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## Probing the impact of metallicity on the dust properties in galaxies

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## Abstract

The aim of this work is to study dust in the interstellar medium (ISM) of nearby galaxies and especially how the metallicity impacts the physical properties of the dust. Local dwarf galaxies, characterised by a low metallicity, are thus the ideal targets for such a study. Indeed previous works have shown that the spectral energy distribution (SED) of these galaxies was different from those of more metal-rich environments. However these studies were limited to the warmer dust emitting at wavelengths shorter than $200 \mu \mathrm{~m}$ and were done only on a limited number of dwarf galaxies. Thanks to its increased sensitivity and resolution in the far-infrared (FIR), Herschel gives us a new view on the cold dust properties in galaxies and enables us to study the lowest metallicity galaxies in a systematic way. Our work is focused on a sample of 48 low-metallicity local dwarf galaxies from one of the Herschel key programs: the Dwarf Galaxy Survey (DGS). We complete the DGS with another Herschel sample : Key Insights on Nearby Galaxies : a Far-Infrared Survey with Herschel (KINGFISH), probing more metal-rich environments.

We adopt a multi-wavelength approach, first focusing on the cold dust probed by Herschel in the FIR and second by probing the observed SED over the whole infrared range (using ancillary data from 2MASS, Spitzer, WISE, IRAS and ground-based facilities), that we model with a semiempirical dust model. We find that a good wavelength coverage of the peak of the SED and a realistic dust model is crucial to properly determine the dust properties, especially the dust mass. We find that dust masses estimated using a modified blackbody model are systematically underestimated by a factor of $\sim 1.8$ compared to dust masses from a full dust SED model. We find that dwarf galaxies harbour on average warmer dust than their metal-rich counterparts (about 1.5 times warmer), due to the harder interstellar radiation field interacting with the porous ISM of dwarf galaxies.

Previous observations of some low-metallicity dwarf galaxies also led to the detection of an emission in the submillimeter (submm) that current models cannot explain: the so-called "submm excess". In our sample of 110 galaxies we confirm the presence of a submm excess in half of the sample, and not necessarily metal-poor objects. We find that submm colours are a good empirical tool to detect or predict the appearance of the submm excess. Another dust composition, using more emissive grains, is able to explain the excess for half of the galaxies. However another mechanism is needed to account for the strongest excesses.

We finally look at the variation of the gas-to-dust mass ratios (G/D) in our sample with metallicity as it is a good tracer of the chemical evolution stage of a galaxy. In the past, the G/D has been seen to correlate linearily with metallicity down to metallicities of $\sim 1 / 5 Z_{\odot}$. We add another sample to the DGS and KINGFISH samples, totalling 126 galaxies, spanning a 2 dex metallicity range, from $\sim 1 / 40 Z_{\odot}$ to $\sim 2 Z_{\odot}$. Our study shows that this linear dependance does not hold at low metallicities and that the relation between $G / D$ and metallicity is steeper. We provide an empirical relation, via a broken power-law fit to the observed G/D values, to estimate the G/D at low metallicities from a metallicity value. However, the scatter in the observed G/D vs metallicity relation is large, due to the variety of environments that we probe, showing that metallicity is not the only parameter driving the relation. We find that the star-formation histories of the galaxies and dust grain growth in the ISM are both important in the evolution of the G/D with metallicity.

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## Introduction

The work presented in this manuscript is incorporated within the vast and general framework of evolution of galaxies. The processes by which galaxies evolve, from the very first galaxies of the primordial Universe to the present-day systems, are still widely debated. The seeds of this evolution lay in the star-formation histories of the galaxies. How are stars formed and at what rate? What controls the efficiency at which the stars are formed? How does this efficiency vary over time? How does this impact the galaxy?

Galaxies are composed primarily of stars and of an interstellar medium (ISM) of gas and dust. The continuous interplay between stars and the ISM is one of the major drivers of galaxy evolution. The ISM plays a key role in this evolution, being the repository of stellar ejecta and the site of stellar birth. It thus contains the imprint of the astrophysical processes occurring in a galaxy. These astrophysical processes can be studied with observations of ISM in galaxies at various wavelengths throughout the spectral energy distribution (SED). The SED is a powerful tool to probe the current physical and chemical state of a galaxy with this historical information integrated over the lifetime of the galaxy.

Interstellar dust is present in most phases of the ISM, from warm ionized regions around young stars to the cores of dense molecular clouds. Because dust is mainly formed from the available metals in the ISM, the dust content traces its internal evolution through metal enrichment. Metallicity is thus a key parameter in studying the evolution of galaxies and is at the heart of this work. Dust influences the subsequent star formation and has a significant impact on the total SED of a galaxy: the absorbed stellar light by dust in the ultraviolet (UV) and visible wavelengths is re-emitted in the infrared (IR) domain by the dust grains. In our Galaxy, dust reprocesses about $30 \%$ of the stellar power, and it can be as large as $\sim 99 \%$ in a starburst galaxy. Studying the IR emission of galaxies thus provides valuable information on the dust properties of the galaxies and on their overall star-formation activity.

With the advent of infrared space missions, such as ISO, IRAS, Spitzer, AKARI, and more recently Herschel, much progress has been made in characterizing the dust properties of galaxies. For a few decades now, IR astronomy has revolutionized our understanding of the dust properties in galaxies. Dust is no longer viewed only as the culprit in obscuring starlight, but also as a powerful tracer of the star-formation activity and of the evolutionary stage of the galaxy. Our Galaxy, as well as other well studied local Universe galaxies, provide various observational benchmarks to calibrate the physical dust properties around solar metallicity. However, for galaxies of the high-redshift Universe, dust properties are still poorly known, and evolution of the dust properties with galactic properties, such as metallicity, star-formation activity, morphology, etc. is unclear. Because of their low metal abundance and active star formation, dwarf galaxies of the local Universe are ideal laboratories for studying star formation and its feedback on the ISM in conditions that may be representative of different stages in early Universe environments.

However due to their low luminosities, and to the lack of full wavelength coverage from the mid-IR to submillimetre (submm) wavelengths, only a handful of dwarf galaxies had been studied until recently, and large uncertainties existed on their dust properties, even on one as fundamental
as the dust mass. An excess emission in the submm has also been observed in several dwarf galaxies and low-metallicity spirals, unexplained by current dust models. With the new wavelength range covered by Herschel in the far-IR (FIR) and submm domains and its unprecedented sensitivity, we are now able to study the detailed dust properties more accurately, and probe dust in systems of extremely low metallicities ( $\sim 1 / 40 Z_{\odot}$ ).

My research focus on the following questions: How do the dust properties (such as mass, temperature, luminosity, etc.) vary with metallicity? In which environments does the submm excess preferentially appear? How do the proportions of metals in the gas and dust phases relate to each other? How do the gas-to-dust mass ratios (G/D) evolve as a function of metallicity?

In this work, I carry out a study of the dust properties in dwarf galaxies and compare with more metal rich environments, in order to address the question of the impact of metallicity on the dust properties. The novelty of this work lays in the fact that dwarf galaxies are studied here in a systematic way, enabling us to derive and quantify the general properties that are representative of these systems. This study is conducted over the full IR-to-submm range, using new FIR/submm Herschel observations, along with Spitzer, WISE, IRAS and 2MASS data. We complete this set of data with longer submm measurements from ground-based facilities such as APEX and JCMT to study the presence and characteristics of the submm excess in my sample of galaxies. I also collect Hi and CO data to access the gas properties of the galaxies and study the evolution of the G/D with metallicity.

This manuscript is composed of three parts. The first part is dedicated to a general description of the ISM in galaxies (Chapter 1) and of low-metallicity systems and their peculiarities (Chapter 2). In these first two Chapters, we also present the main motivations for this work. The second part of my work is dedicated to the study of the cold dust component uncovered by Herschel in lowmetallicity environments. Herschel and the two samples that are considered for this study (a sample of dwarf galaxies and a sample probing more metal-rich environments) are presented in Chapter 3. Chapter 4 presents the new Herschel data obtained for our objects, that we interpret with the aid of a modified blackbody model (Chapter 5). The third part of my work (Chapters 6 to 9 ) broadens our view of dust in these low-metallicity systems by looking at the total dust component across the whole IR-to-submm range (data presented in Chatper 6). The SEDs are modelled with a semiempirical dust model, presented in Chapter 7 and the results are then analysed and discussed in Chapter 8. We especially focus on the variation of the dust properties with metallicity and address the submm excess issue in this Chapter. The last chapter of this thesis (Chapter 9) incorporates the gas into our dust picture and presents a study of the variations of G/D of galaxies with metallicity. Results are discussed and interpreted with the aid of three chemical evolution models.

## Part I

## The ISM of low-metallicity galaxies

## Chapter 1

## Structure and physics of the interstellar medium

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This Chapter aims to give a brief overview of our current understanding of the structure and the physics of the interstellar medium (ISM) of galaxies and especially our Galaxy for which most details are known. Our current view of the ISM has been shaped by the observations at various wavelengths available to astronomers. Thus we start this Chapter with a description of the physical properties that can be probed by observations at different wavelengths (Section 1.1), and introduce the main components of a galaxy. Then, detailed properties of the dust and gas, the two main components of the ISM, are presented and discussed in Sections 1.2 and 1.3 for the case of our Galaxy. We especially focus on linking the observations of the ISM with a theoretical description of the physical processes at stake responsible for what we see. In the last Section 1.4 we describe the main tools at our disposal to probe the ISM emission of a galaxy, and describe how this emission can vary according to the different types of galaxies.

### 1.1 A multi-wavelength picture of a galaxy

In the most general view, a galaxy is composed of stars evolving in an interstellar medium (ISM) and surrounded by a dark matter halo. The ISM contains all of the matter that fills in the vast empty space between the stars in a galaxy: for example, in the solar neighbourhood, stars are separated by $\sim 2 \mathrm{pc}$. The ISM can be decomposed into a gaseous component: interstellar gas; and a solid component: interstellar dust. Gas structures the ISM in a large variety of phases depending on its local physical conditions (mainly density and temperature): from ionized to molecular phase, where "ionized" and "molecular" refer to the state of the hydrogen element. Dust, however, is present throughout the ISM.

The only way to access information about a galaxy, such as its composition or its physical and chemical states, is through the light we receive from space. Galaxies emit over the whole electromagnetic spectrum and the light at a given energy brings its own piece of information: i.e., having access to a large range of wavelengths is crucial to build the most complete picture of a given galaxy. Fig. 1.1 shows, as an example, the Whirlpool galaxy, M51, seen with five different instruments ${ }^{1}$.


Fig. 1.1. M51 seen by Chandra (hot gas), GALEX (young stars), HST (old stars), Spitzer (dust) and VLA (atomic gas) telescopes.

We clearly see differences between the various wavelengths: for example the spiral structure of the galaxy is visible at all the wavelengths observed here except in the X-rays; or the northern companion of M51 is not visible with GALEX. We can deduce the presence of all of the components in the galaxy with these observations: for example, stars are visible with GALEX and HST. If there are stars, then gas must also be present as can be seen with Chandra and the VLA. And finally the

[^0]black lanes obscuring the HST image can be attributed to micro-particles obscuring the light from the galaxy, that were generically called dust.

The distribution of the amount of total energy emitted by a galaxy over a given wavelength or frequency range is called a spectral energy distribution (SED). The SED is a precious tool for the astronomers as it gives an instantaneous multiwavelength snapshot of a galaxy. Figure 1.2 shows an example of UV-to-mm SED for a galaxy. The black points and red spectrum represent direct observations from the galaxy. The challenge for astronomers is to find a realistic model (i.e., the blue curve of Fig. 1.2) that will take into account all of the physical processes occurring in the galaxy, in order to disentangle the emission from the different components and then access information about the galaxy. Here for example, the model tells us that this galaxy is forming stars at a rate of 9.5 $\mathrm{M}_{\odot} \mathrm{yr}^{-1}$.


Fig. 1.2. Example of a UV-to-mm SED for a galaxy, here NGC7714 from Groves et al. (2008).The black points and red spectrum are observational constraints. A realistic model (blue curve from Groves et al. 2008) reproduces the emission from the galaxy and enables us to access physical properties of the galaxy. Note: We define the acronyms for the wavelength ranges we use throughout this manuscript: ultra-violet (UV) from 0.1 to $0.4 \mu \mathrm{~m}$, visible from 0.4 to $0.8 \mu \mathrm{~m}$, near-infrared (NIR) from 0.8 to $\sim 5 \mu \mathrm{~m}$, mid-infrared (MIR) from $\sim 5$ to $\sim 50 \mu \mathrm{~m}$, far-infrared (FIR) from $\sim 60$ to $\sim 600-700 \mu \mathrm{~m}$ and sub-millimetre (submm) from $\sim 500$ to $1000 \mu \mathrm{~m}$.

Indeed the various components of a galaxy all dominate the total emission in a specific wavelength domain:

Stars: Stars dominate in the UV to NIR range: young stars (ages of a few Myr) have a peak of emission around $0.1 \mu \mathrm{~m}$ in the UV and older stars (ages of $0.1-1 \mathrm{Gyr}$ ) have their emission peaking in the visible around $0.5 \mu \mathrm{~m}$. Thus in Fig. 1.2, stars are responsible for the UV-to-NIR peak of the SED.

Dust: Dust absorbs the stellar radiation and re-emits the bulk of its energy in the IR. Emission by dust will be discussed in Section 1.2.4 and will be the main topic of Chapter 7. The broad IR peak of the SED in Fig. 1.2 is due to dust in NGC7714.

Gas: Gas emits over the whole electromagnetic spectrum and is visible in the total SED of a galaxy through spectral lines and a continuum in emission or in absorption. These spectral lines come from a change in the state of the gas (excitation) and are powerful tracers of the physical state of the gas in the galaxy (see Section 1.3). Atomic hydrogen gas emission can dominate the SED of a galaxy in radio wavelengths.

We note in Fig. 1.2 a sudden and important drop in the energy emitted at far-UV (FUV) wavelengths around $0.1 \mu \mathrm{~m}$. This is because, at wavelengths below the Lyman edge at $912 \AA$, a small amount of hydrogen present in the ISM is capable of absorbing nearly all radiation. This sharp transition at $912 \AA$ is often used by cosmologists to determine the redshift of a galaxy: knowing the wavelength of this transition in the observed spectrum of a distant galaxy and comparing to $912 \AA$ indicates its redshift. At shorter wavelengths in the X-ray emission domain, the total emission of a galaxy is dominated by emission from hot plasmas: coronal gas in the halo of the galaxy and in supernovae remnants.

The relative importance of all of these components (stars, gas and dust) to the total SED is not the same for different galaxies. Indeed these differences depend on numerous physical characteristics of the galaxies such as morphological type or stellar activity for example (see Section 1.4). To be able to disentangle the different components, we need to know more about how they interact with each other, about the physical processes occurring in the galaxy and the theory governing them. This will be the object of Sections 1.2 and 1.3.

Let us now focus more in details on the ISM of our Galaxy. Table 1.1 gives a view of the repartition of mass and luminosity between the different components of the Galaxy, and provides useful reference values for comparison with other galaxies. The total baryonic mass is dominated by the mass of stars and the ISM represents about $2.5 \%$ of the baryonic mass (Tielens 2005), but it accounts for about $30 \%$ of the total luminosity.

Table 1.1. Masses and luminosities in the Milky Way (adapted from Table 3 of Tielens 2005).

| Component | Mass $\left[M_{\odot}\right]$ | Luminosity $\left[L_{\odot}\right]$ |
| :--- | :---: | :---: |
| Stars | $1.8 \times 10^{11}$ | $\sim 4.8 \times 10^{10}$ |
| Gas | $4.5 \times 10^{9}$ | $\sim 2 \times 10^{8}$ |
| Dust | $\sim 4.5 \times 10^{7}$ | $\sim 1.7 \times 10^{10}$ |
| Total | $\sim 2 \times 10^{11, a}-\geq 10^{12, b}$ | $\sim 6.5 \times 10^{10}$ |

${ }^{a}$ : visible matter only.
${ }^{b}$ : including dark matter.

### 1.2 Dust in the ISM

### 1.2.1 Observational evidence of dust in the ISM

The first evidence of the presence of dust in the ISM was through the obscuration of some parts of the sky in the optical wavelengths that were first accessible to astronomers (Fig. 1.3 ${ }^{2}$ ). The light coming from the stars is extinguished by dust present along the line-of-sight between the stars and the observer. Comparing the light arriving from a non-obscured star and that from an obscured star, i.e., measuring extinction (see Section 1.2.2), provides valuable information on the nature of dust present along the line-of-sight.

[^1]

Fig. 1.3. Optical image of the Galactic Center. Dust present in between the observer and the center of the Milky Way causes the light coming from the inner parts of the Galaxy to be heavily extinguished.

The physical processes responsible for the extinction of stellar light are absorption and scattering of the light by dust grains, which strongly depend on the intrinsic properties of the dust grain, mainly its chemical composition and size. We observe extinction of starlight along a wide range of wavelengths indicating that grains of different natures and sizes are present in the ISM. A dust model can be defined as an ensemble of dust properties (see Section 1.2.5) needed to reproduce a set of observational constraints. That is why the very first dust models (e.g., Mathis et al. 1977) were derived using the mean interstellar extinction in our Galaxy to constrain the dust size distribution and chemical composition.

As the dust grain absorbs energy, it also radiates this energy away at IR wavelengths. Thus we can also obtain constraints on the nature of the dust by looking at emission of the ISM in the IR domain. This was only possible after the advent of space astronomy in the early 80 's with the first IR space telescope: the InfraRed Astronomy Satellite (IRAS). Indeed the Earth atmosphere absorbs IR radiation from space and obliges the astronomers to explore this wavelength range from space. The availability of these new constraints, from dust emission, revolutionized the study of ISM dust, as astronomers started to incorpore them into their models (e.g., Desert et al. 1990).

Dust grains are made from the available elements in the ISM, the most abundant being carbon and oxygen. A fraction of them being locked up in the solid phase of the ISM, these elements are depleted from the gas phase. Informations about the chemical composition of the dust can thus be deduced from what is called "elemental depletions", i.e., what is not seen in the gas phase (as in the modelling of Zubko et al. 2004).

Other constraints can be used to determine the nature of the dust grains such as X-ray scattering by dust (Smith \& Dwek 1998), bringing information on the abundance and composition of the biggest grains; polarisation of the starlight by dust providing information on the geometry of the medium and on the grain elongation; reflection nebulae constraining the grain albedo ${ }^{3}$; or spectral features to access the grain composition.

[^2]
### 1.2.2 Extinction by dust grains

Dust has thus a major impact on the propagation of starlight throughout the galaxy as it absorbs radiation from the stars. The amount of extinction suffered from a starlight radiation after passing through a cloud of dust is related to the optical depth of the cloud $\tau(\lambda)$, which is, in turn, related to the dust column density, i.e., the quantity of dust along a given line-of-sight. Historically, the extinction is expressed as $A(\lambda)$, in magnitudes (see Eq. 1.2).

The monochromatic specific intensity received by the observer, $I_{\lambda}$, after passing through a cloud of dust with an optical depth $\tau(\lambda)$, is related to the source intensity, $I_{\lambda}^{0}$, by:

$$
\begin{equation*}
I_{\lambda}=I_{\lambda}^{0} \times e^{-\tau(\lambda)}=I_{\lambda}^{0} \times 10^{-A(\lambda) / 2.5} \tag{1.1}
\end{equation*}
$$

From Eq. 1.1 we can relate $\tau(\lambda)$ and $A(\lambda)$ :

$$
\begin{equation*}
A(\lambda)=1.086 \times \tau(\lambda) \tag{1.2}
\end{equation*}
$$

Figure 1.4 shows the extinction curve of the diffuse Galactic ISM from UV to mm wavelengths. Note the two absorption features at 9.7 and $18 \mu \mathrm{~m}$ and the "UV-bump" at $2175 \AA$, that will be attributed, respectively, to silicates, and small graphitic grains and large molecules (see Section $1.2 .5)$. We see that the extinction is most important in the UV and optical wavelengths and drops at longer wavelengths.


Fig. 1.4. Extinction curve for the diffuse Galactic ISM, from Mathis (1990) with the silicate features from Dudley \& Wynn-Williams (1997). The $\lambda^{-2}$ dependence in the IR is indicated with the grey dashed line (see Section 1.2.3).

By definition $A(\lambda)$ depends on where the observer is looking in the sky: the form of the extinction curve is dictated by the grain properties and this can differ depending on the local environment. Thus the extinction curve can vary across the Galaxy and from one galaxy to another (see Chapter $2)$.

## Extinction determination

The most common way of estimating $A(\lambda)$ is via the "pair-method". It consists in comparing the SEDs of two stars of the same spectral type and luminosity class, i.e., two stars with the same absolute magnitude, Mag. The "reference" star is assumed to be unaffected by extinction and the other star is located behind the absorbing material for which we want to know $A(\lambda)$. At a given wavelength their apparent magnitudes, mag, are given by:

$$
\left\{\begin{align*}
\operatorname{mag}(\lambda) & =\operatorname{Mag}(\lambda)+5 \times \log (d)+A(\lambda)  \tag{1.3}\\
\operatorname{mag}_{r e f}(\lambda) & =\operatorname{Mag}_{\text {ref }}(\lambda)+5 \times \log \left(d_{r e f}\right)
\end{align*}\right.
$$

where $d$ and $d_{\text {ref }}$ are the distances to the star and to the "reference" star respectively. Given that $\operatorname{Mag}(\lambda)=\operatorname{Mag}_{r e f}(\lambda)$, we can access $A(\lambda)$ by comparing the apparent magnitudes at two different wavelengths, $\lambda$ and $\lambda_{0}$. By taking the difference of the apparent magnitudes, we remove the unknown distance parameter:

$$
\begin{equation*}
A(\lambda)-A\left(\lambda_{0}\right)=\left[\operatorname{mag}(\lambda)-\operatorname{mag}_{r e f}(\lambda)\right]-\left[\operatorname{mag}\left(\lambda_{0}\right)-\operatorname{mag}_{r e f}\left(\lambda_{0}\right)\right] \tag{1.4}
\end{equation*}
$$

or:

$$
\begin{equation*}
A(\lambda)-A\left(\lambda_{0}\right)=\Delta \operatorname{mag}(\lambda)-\Delta \operatorname{mag}\left(\lambda_{0}\right) \tag{1.5}
\end{equation*}
$$

The colour excess between two wavelengths, $E\left(\lambda-\lambda_{0}\right)$, is defined by $\Delta \operatorname{mag}(\lambda)-\Delta \operatorname{mag}\left(\lambda_{0}\right)$ and usually the extinction is expressed in terms of $E(\lambda-V)$ and normalised by $E(B-V)=A_{B}-A_{V}$, where B and V refer to the optical bands at 440 and 560 nm . An important parameter can be extracted from the extinction curve in Fig. 1.4: $R_{V}=A_{V} / E(B-V)$. This parameter describes the ratio of the total-to-selective extinction and can be estimated from the IR slope of the extinction curve. In the diffuse Galactic ISM, the mean value of $R_{V}$ is $3.05 \pm 0.15$ (Whittet 2003). The value of $R_{V}$ depends on the nature and on the size distribution of the dust grains along the line-of-sight. For example, in denser regions, where bigger grains can be formed by coagulation, $R_{V}$ is higher $(\sim 5)$.

### 1.2.3 Dust physics: absorption and scattering processes

As we saw in Section 1.2.1, dust extinguishes stellar light. The physical processes responsible for this extinction are absorption and scattering of the light by dust grains. These processes depend strongly on the intrinsic properties of the dust grains, as described in this Section.

## Useful definitions and notations

Let us now consider a homogeneous cloud of spherical dust grains of radius $a$. The dust optical depth, $\tau(\lambda)$, through a dust cloud of size $l$, is related to the dust properties via:

$$
\begin{equation*}
\tau(\lambda)=n_{d} C_{e x t}(a, \lambda) l \tag{1.6}
\end{equation*}
$$

where $n_{d}$ is the dust number density, $C_{\text {ext }}$ is the dust extinction cross-section at wavelength $\lambda$ for a grain radius $a$.

For a dust grain, $C_{\text {ext }}$ is given by:

$$
\begin{equation*}
C_{e x t}(a, \lambda)=Q_{e x t}(a, \lambda) \times \pi a^{2} \tag{1.7}
\end{equation*}
$$

where $\pi a^{2}$ represents the geometric cross-section of the dust grain and $Q_{\text {ext }}$ is the extinction efficiency. As extinction is due to absorption and scattering, $C_{\text {ext }}$ can be decomposed into:

$$
\begin{equation*}
C_{e x t}=C_{a b s}+C_{s c a} \tag{1.8}
\end{equation*}
$$

with:

$$
\left\{\begin{array}{l}
C_{a b s}(a, \lambda)=Q_{a b s}(a, \lambda) \times \pi a^{2}  \tag{1.9}\\
C_{s c a}(a, \lambda)=Q_{s c a}(a, \lambda) \times \pi a^{2}
\end{array}\right.
$$

where $Q_{a b s}$ and $Q_{s c a}$ are the absorption and scattering efficiencies respectively.
We can define the "absorption coefficient", $\alpha(\lambda)$, of the dust cloud by:

$$
\begin{equation*}
\alpha(\lambda)=n_{d} C_{a b s}(a, \lambda) \tag{1.10}
\end{equation*}
$$

Often, $\alpha(\lambda)$ is written as $\alpha(\lambda)=\rho_{I S M} \kappa_{a b s}(\lambda)$, with $\rho_{I S M}$ the cloud mass density and $\kappa_{a b s}(\lambda)$ is called the "opacity" or "mass absorption coefficient". $\kappa_{a b s}(\lambda)$ is given by:

$$
\begin{equation*}
\kappa_{a b s}(\lambda)=\frac{n_{d} \pi a^{2} Q_{a b s}(a, \lambda)}{\rho_{I S M}} \tag{1.11}
\end{equation*}
$$

We have:

$$
\begin{equation*}
\frac{n_{d}}{\rho_{I S M}}=\frac{N_{g}}{V_{I S M}} \frac{V_{I S M}}{M_{g}}=\frac{1}{m_{g}} \tag{1.12}
\end{equation*}
$$

where $N_{g}$ is the number of dust grains along the line-of-sight within the beam, $M_{g}$ is the total mass of the grains, and $m_{g}$ is the mass of a single dust grain: $m_{g}=\frac{4}{3} \pi a^{3} \rho$, where $\rho$ is the dust material mass density. Injecting Eq. 1.12 into Eq. 1.11 yields:

$$
\begin{equation*}
\kappa_{a b s}(\lambda)=\frac{\pi a^{2} Q_{a b s}(a, \lambda)}{m_{g}}=\frac{3 Q_{a b s}(a, \lambda)}{4 a \rho} \tag{1.13}
\end{equation*}
$$

## Accessing the dust properties: absorption and scattering efficiencies

The quantities $Q_{a b s}$ and $Q_{s c a}$ contain all of the dust grain information: its chemical components, its porosity and the spectral dependence of its dielectric properties. They represent an intrinsic property of the dust grain. These quantities can be measured in the laboratory or theoretically derived by solving Maxwell's equations of propagation of an electromagnetic wave through the dust grain with the appropriate boundary conditions. In this theoretical framework, the dust grain is made of a material with complex refractory index, $m$, related to the dielectric constant, $e_{\text {diel }}$, of the material via $e_{\text {diel }}=m^{2}$. The Mie theory provides numerical solutions for absorption and scattering in the case of a spherical and homogeneous dust grain. This theory has also been extended to more general cases of non-spherical dust grains. For example, realistic dielectric functions for graphite and silicates have been derived by Draine \& Lee (1984) (see Section 1.2.5 for reasons to consider these specific materials in the interpretation of observations).

In the case of spherical dust grains, the absorption and scattering efficiencies can be expressed as series expansions of the "size parameter", $x=2 \pi a / \lambda$. We can distinguish two limiting regimes:

The grey extinction: $\mathrm{x} \gg 1$ i.e., the wavelength is very small compared to the grain size and the grain behaves as an opaque screen. We have:

$$
\left\{\begin{array}{l}
Q_{a b s}(a, \lambda) \sim 1  \tag{1.14}\\
Q_{s c a}(a, \lambda) \sim 1
\end{array}\right.
$$

The Rayleigh limit: $\mathrm{x} \ll 1$ i.e., the grain size is very small compared to the wavelength, as it is the
case in the IR. In this domain, only the first term of the series dominates and we have:

$$
\begin{equation*}
Q_{a b s}(a, \lambda)=4 x \operatorname{Im}\left[\frac{m^{2}-1}{m^{2}+2}\right] \tag{1.15}
\end{equation*}
$$

where Im denotes the imaginary part of the bracketed expression, and:

$$
\begin{equation*}
Q_{s c a}(a, \lambda)=\frac{8}{3} x^{4}\left|\frac{m^{2}-1}{m^{2}+2}\right| \tag{1.16}
\end{equation*}
$$

Often, $Q_{a b s}$ is approximated in the IR domain by a power-law with index $\beta$ :

$$
\begin{equation*}
Q_{a b s}(a, \lambda)=Q_{0}\left(\frac{\lambda_{0}}{\lambda}\right)^{\beta} \text { for } \lambda>\lambda_{0} \tag{1.17}
\end{equation*}
$$

where $\beta$ is called the "emissivity index", and $Q_{0}$ is such that $Q_{a b s}\left(a, \lambda_{0}\right)=Q_{0}$. Typically, values of $\beta$ between 1 and 2 are considered. Here we develop why that is.

Let us assume that we can model a dust grain by a system of harmonic oscillators, representing the plasma of free electrons tied to the lattice of the crystal, and that this dust grain is interacting with a plane electromagnetic wave with electric field $E=E_{0} e^{i \omega t}$, where $\omega$ is the pulsation ( $\omega=2 \pi \nu$ ). After solving the equations of motion for a harmonic oscillator we can derive a dispersion relation and obtain an expression of the dielectric function, $e_{\text {diel }}$, as a function of the pulsation, $\omega$ :

$$
\begin{equation*}
e_{\text {diel }}=1+\frac{\omega_{p}^{2}\left(\omega_{0}^{2}-\omega^{2}\right)}{\left(\omega_{0}^{2}-\omega^{2}\right)^{2}+\gamma_{0}^{2} \omega^{2}}+i \frac{\omega \omega_{p}^{2} \gamma_{0}}{\left(\omega_{0}^{2}-\omega^{2}\right)^{2}+\gamma_{0}^{2} \omega^{2}} \tag{1.18}
\end{equation*}
$$

where $\omega_{0}$ and $\gamma_{0}$ are the resonant frequency and the damping constant, respectively, of the oscillator; and $\omega_{p}$ is the plasma pulsation: $\omega_{p}=\sqrt{\frac{4 \pi N e^{2}}{m_{e}}}$, where $N$ is the number density of free electrons.

In the ISM, we can assume that the electromagnetic field interacting with the grain is slowly varying. In this approximation: $\omega \ll \omega_{0}$, and the expression for $e_{\text {diel }}$ becomes:

$$
\begin{equation*}
e_{\text {diel }}=1+\frac{\omega_{p}^{2}}{\omega_{0}^{2}}+i \frac{\gamma_{0} \omega_{p}^{2}}{\omega_{0}^{4}} \omega \tag{1.19}
\end{equation*}
$$

Using $e_{\text {diel }}=m^{2}$ and writing $e_{\text {diel }}=e_{1}+\mathrm{i} e_{2}$, we can transform Eq. 1.15 in:

$$
\begin{equation*}
Q_{a b s}(a, \lambda)=\frac{8 \pi a}{\lambda} \frac{3 e_{2}}{\left|e_{d i e l}+2\right|^{2}} \tag{1.20}
\end{equation*}
$$

Injecting Eq. 1.19 in Eq 1.20 gives:

$$
\begin{equation*}
Q_{a b s}(a, \lambda)=\frac{48 \pi^{2} a c}{\lambda^{2}} \frac{\gamma_{0} \omega_{p}^{2}}{\omega_{0}^{4}\left(3+\frac{\omega_{p}^{2}}{\omega_{0}^{2}}\right)^{2}} \tag{1.21}
\end{equation*}
$$

Thus for an idealised dust grain made of harmonic oscillators, in the slowly varying field approximation and Rayleigh limit we have:

$$
\begin{equation*}
Q_{a b s}(a, \lambda) \sim \lambda^{-2} \tag{1.22}
\end{equation*}
$$

Thus for a crystalline grain we have $\beta=2$. In a more general case, materials can show a range of $\beta$ values depending on their temperature.

Following the same method for the scattering efficiency we also get:

$$
\begin{equation*}
Q_{s c a}(a, \lambda)=\frac{8}{3}\left(\frac{2 \pi a}{\lambda}\right)^{4} \frac{\omega_{p}^{2}}{\omega_{0}^{2}\left(3+\frac{\omega_{p}^{2}}{\omega_{0}^{2}}\right)} \tag{1.23}
\end{equation*}
$$

i.e.,

$$
\begin{equation*}
Q_{s c a}(a, \lambda) \sim \lambda^{-4} \tag{1.24}
\end{equation*}
$$

From Eqs. 1.22 and 1.24, we see that at IR wavelengths, absorption dominates over scattering.


Fig. 1.5. Extinction efficiencies, $Q_{e x t}=Q_{a b s}+Q_{s c a}$, from 0.01 to $1000 \mu \mathrm{~m}$, for silicate grains (in blue, from Weingartner \& Draine 2001) and graphite grains (in green, from Laor \& Draine 1993) of different radiuses: $a=0.001 \mu \mathrm{~m}$ (dotted lines), $0.01 \mu \mathrm{~m}$ (plain lines) and $0.1 \mu \mathrm{~m}$ (dashed lines). The $\lambda^{-2}$ dependence in the IR is indicated with the grey dashed line.

Figure 1.5 shows the extinction efficiency, $Q_{e x t}$, for different grains of different sizes. We see that for big grains the extinction efficiency is constant at short wavelengths (UV - visible): we are in the "grey extinction" limit (Eq. 1.14). In the IR domain, the grain size is small compared to the wavelength and $Q_{e x t} \sim Q_{a b s} \sim \lambda^{-2}$ (Rayleigh limit, Eq. 1.15 and 1.22). As $Q_{e x t}$ is directly linked to $\mathrm{A}(\lambda)$ through Eqs. 1.2, 1.6, and 1.7, we find the same $\lambda^{-2}$ dependency in the FIR for the extinction curve, as presented in Fig. 1.4. We also note that, for a given grain size, in the UV - visible the graphite extinction efficiency is greater than the silicate extinction efficiency. The opposite is true in the MIR. Both types are equivalent in the FIR. We clearly see here the influence of the nature of the grain on the observed dust properties.

### 1.2.4 Emission by dust grains

After absorbing stellar light, the dust grains re-emit the absorbed energy and this emission depends on the nature of the grains and on the strength and hardness of the incident radiation field.

## The interstellar radiation field

The ISM of a galaxy bathes in the interstellar radiation field (ISRF) created by every source of radiation in the galaxy. Three sources of radiation can be identified:
Stars: They are the main source of radiation in the ISM, and play a prominent role in determining the thermal balance of the dust grains.
Gas and Dust: Gas and dust are not a priori radiation sources. They radiate away the energy absorbed from stars, in the IR range for the dust and over a large part of the electromagnetic spectrum for the gas.
Extragalactic component: It is mainly composed of the cosmic microwave background (CMB) radiation in the millimetre ( mm ) wavelength range. The CMB is a radiation of cosmological origin with the spectral shape of a $T_{C M B}=2.73 \mathrm{~K}$ blackbody (Mather et al. 1994). This is the only radiation, along with cosmic rays, that can penetrate into deeply embedded regions. This means that everywhere in the ISM, even in this deeply embedded regions, dust grains have a minimal temperature equal to $T_{C M B}$.

Technically all of the three sources of radiation contribute to dust heating. However, only the stellar component provides photons energetic enough to sensibly affect the internal energy of the grains and to induce a significant change in temperature (except in the deepest embedded regions where the CMB and cosmic rays also contribute to the dust heating). Thus we will only focus on the UV-to-NIR part of the ISRF.

In the solar neighbourhood, the ISRF has been empirically determined by Mathis et al. (1983), and is composed of: a UV component to model the young stars; three blackbodies with different temperatures $T_{\{1,2,3\}}=\{7500,4000,3000\} \mathrm{K}$, to model the main disk stellar populations and old stars; and the CMB radiation. A zoom on the UV-to-NIR part of the Mathis et al. (1983) ISRF is shown in Fig. 1.6.

## Grains in thermal equilibrium

Inside the dust cloud, a dust grain large enough and bathing in a given interstellar radiation field acquires an equilibrium temperature, $T_{d}$. This temperature is determined by the condition that the dust grain absorbs the same amount of energy as it emits per unit time: i.e., the power absorbed is the same as the power emitted.

The power (or luminosity) absorbed by the dust cloud is given by:

$$
\begin{equation*}
\Gamma_{a b s}=\int 4 \pi \alpha(\lambda) J_{\lambda}(\lambda) \mathrm{d} \lambda=\int 4 \pi \times n_{d} \pi a^{2} Q_{a b s}(a, \lambda) J_{\lambda}(\lambda) \mathrm{d} \lambda \tag{1.25}
\end{equation*}
$$

where $J_{\lambda}$ is the mean intensity of the interstellar radiation field, and $\alpha(\lambda)$ is the absorption coefficient from Eq. 1.10.

The power (or luminosity) emitted by the dust cloud is given by:

$$
\begin{equation*}
\Gamma_{e m}=\int 4 \pi j_{\lambda}(\lambda) \mathrm{d} \lambda \tag{1.26}
\end{equation*}
$$

where $j_{\lambda}$ is the emission coefficient, in units of $\mathrm{W} \mathrm{m}{ }^{-3} \mathrm{sr}^{-1} \mu \mathrm{~m}^{-1}$.
Kirchhoff's law for thermal emission tells us that $j_{\lambda}=\alpha(\lambda) \times B_{\lambda}\left(\lambda, T_{d}\right)$ where $B_{\lambda}(\lambda, T)$ is the Planck function. Injecting in Eq. 1.26, we have:

$$
\begin{equation*}
\Gamma_{e m}=\int 4 \pi \times n_{d} \pi a^{2} Q_{a b s}(a, \lambda) B_{\lambda}\left(\lambda, T_{d}\right) \mathrm{d} \lambda \tag{1.27}
\end{equation*}
$$

At thermal equilibrium we have $\Gamma_{a b s}=\Gamma_{e m}$, leading to:


Fig. 1.6. Zoom on the UV-to-NIR part of the Mathis et al. (1983) ISRF. The UV component is shown in purple, and the three effective stellar sources in blue, green and orange. The total ISRF is shown with the black dashed line.

$$
\begin{equation*}
\int Q_{a b s}(a, \lambda) J_{\lambda}(\lambda) \mathrm{d} \lambda=\int Q_{a b s}(a, \lambda) B_{\lambda}\left(\lambda, T_{d}\right) \mathrm{d} \lambda \tag{1.28}
\end{equation*}
$$

The dust equilibrium temperature $T_{d}$ can be deduced from Eq. 1.28.
$T_{d}$ and Stefan's law In the case of grains in thermal equilibrium, and assuming Eq. 1.17 for the absorption efficiency in the IR, Stefan's law tells us that the energy density ${ }^{4}$ of the radiation emitted by the dust grains is $\propto T_{d}{ }^{4+\beta}$. With $\beta$ usually ranging from 1 to 2 for typical interstellar grains, we see that a relatively small increase of the dust grain temperature results in a large increase of the emitted power of the grain (up to orders of magnitude). In other words, hot grains emit much more than cold grains. Moreover, as the grains are in thermal equilibrium with the interstellar radiation field, we have:

$$
\begin{equation*}
u_{I S R F} \propto T_{d}^{4+\beta} \tag{1.29}
\end{equation*}
$$

where $u_{I S R F}$ is the energy density of the interstellar radiation field integrated over the wavelength. Thus knowing the strength of the interstellar radiation field, one can get a rough estimate of the dust grains temperature. Using the equilibrium temperature value from the diffuse Galactic ISM: $T_{d, \odot}=17.5 \mathrm{~K}$ (Boulanger et al. 1996), we estimate:

$$
\begin{equation*}
T_{d}=T_{d, \odot}\left(\frac{u_{I S R F}}{u_{\text {ISRF }, \odot}}\right)^{1 /(4+\beta)} \tag{1.30}
\end{equation*}
$$

[^3]where $u_{I S R F, \odot}$ is the energy density of the diffuse Galactic ISRF, given by Mathis et al. (1983).
$T_{d}$ and Wien's law Another interesting property of the emission of dust grains in thermal equilibrium is that we can have a useful information on the shape of the SED, the position of the peak of emission (in a $\nu L_{\nu}$ representation), using Wien Displacement's law:
\[

$$
\begin{equation*}
\lambda_{\max } T_{d} \sim 3000 \mu \mathrm{~m} \mathrm{~K} \tag{1.31}
\end{equation*}
$$

\]

i.e., for hotter grains, the peak of the SED is moved to shorter wavelengths. For example, using the value of Boulanger et al. (1996), we can deduce that the peak of dust emission for the diffuse Galactic ISM is around $170 \mu \mathrm{~m}$, in the FIR.
$T_{d}$ and $a$ The dust temperature $T_{d}$ also depends on the grain size $a$. Intuitively we can understand that the larger the grains the colder they are. This also means that the bulk of the dust mass resides in cold grains. Using Eq. 1.31 we see that large, cold grains emit at longer wavelengths than small, hot grains.

We emphasize that these relations are only approximations that are only useful to get a rough idea of the dust emission properties. In the ISM, the dust grains are distributed in sizes and thus in temperatures.

## Stochastic heating

When the grains are very small, their temperature fluctuates and we say that they undergo "stochastic heating". For small grains, the internal energy of the grain is comparable to the energy of a single incident photon and thus will be considerably affected by the absorption of a single photon, resulting in a sudden rise in temperature. The grain then cools down to its equilibrium temperature until it absorbs another photon, giving a "spiky" temperature profile (see Fig. 1.7). Inversely, for a big grain, the internal energy is large compared to the energy of an incident photon. Consequently its absorption does not affect the temperature of the grain: big grains are then in thermal equilibrium with a constant temperature profile. Fig. 1.7 illustrates this point for carbonaceous grains heated by the Galactic ISRF.

To compute the emission of a stochastically heated grain, we then need to know its thermal properties in addition to its optical properties, i.e., its specific heat. The specific heat of a dust grain brings information on how much energy is needed to raise the temperature of the dust grain by 1 K . This enables us to compute the temperature of the grain after absorption of a photon of energy $\epsilon_{\gamma}(a)=\mathrm{h} \nu$ via:

$$
\begin{equation*}
\epsilon_{\gamma}(a)=\frac{4}{3} \pi a^{3} \int_{T_{i}}^{T_{f}} C(T) \mathrm{d} T \tag{1.32}
\end{equation*}
$$

where $C(T)$ is the specific heat per unit volume of the grain, $T_{i}$ and $T_{f}$ are the grain temperatures before and after the absorption of a photon of energy $\epsilon_{\gamma}(a)$.

Let us now define $P(a, T) d T$ the probability that a grain of radius $a$ has a temperature between [ $T, T+d T]$, with $P(a, T)$ the temperature probability distribution, verifying:

$$
\begin{equation*}
\int_{0}^{\infty} P(a, T) \mathrm{d} T=1 \tag{1.33}
\end{equation*}
$$

We assume that radiation from a stochastically heated grain obeys Kirchhoff's law at any time, and the emission coefficient now can be expressed as:


Fig. 1.7. Illustration of stochastic heating of small grains: carbonaceous grains heated by the Galactic ISRF (figure from Draine 2003). $a$ is the radius of the grain and $\tau_{a b s}$ the mean time between photon absorptions. The plotted time range corresponds approximately to a day.

$$
\begin{equation*}
j_{\lambda}(a, \lambda)=\alpha(a, \lambda) \times \int_{0}^{\infty} B_{\lambda}(\lambda, T) P(a, T) \mathrm{d} T \tag{1.34}
\end{equation*}
$$

For a big grain, the temperature only oscillates a little around the equilibrium temperature $T_{d}$ and $P(a, T)$ can be approximated by a $\delta$-function $\delta\left(\mathrm{T}-\mathrm{T}_{d}\right)$ : we arrive back at Kirchhoff's law which we used for big grains in thermal equilibrium: $j_{\lambda}=\alpha(\lambda) \times B_{\lambda}\left(\lambda, T_{d}\right)$.

We refer to Guhathakurta \& Draine (1989) for details on the calculation of the temperature probability distribution.

## Grain size distribution

So far we have considered the case of a cloud of dust grains, all with the same radius $a$, absorbing stellar radiation over a large range of wavelengths. In the ISM dust grains exist in a distribution of sizes so let us now consider a dust specie spanning a range in sizes between $a_{\min }$ and $a_{\max }$, the minimum and maximum grain sizes. We define $n(a) d a$ the number density of grains with sizes between $[a, a+d a]$, and $n(a)$ is the size distribution of the dust specie. The dust column density, $N_{d}$, is the number of grains over a given line of sight of length $l$ and is given by:

$$
\begin{equation*}
N_{d}=l \int_{a_{\min }}^{a_{\max }} n(a) \mathrm{d} a \tag{1.35}
\end{equation*}
$$

Following Eq. 1.6, the optical depth, $\tau(\lambda)$, for this dust specie is given by:

$$
\begin{equation*}
\tau(\lambda)=l \int_{a_{\min }}^{a_{\max }} C_{e x t}(a, \lambda) n(a) \mathrm{d} a \tag{1.36}
\end{equation*}
$$

Following Eqs 1.26 and 1.34, the IR luminosity, $L_{\lambda}$, emitted by this dust specie is given by:

$$
\begin{equation*}
L_{\lambda}(\lambda)=l \int_{a_{\min }}^{a_{\max }} 4 \pi^{2} a^{2} Q_{a b s}(a, \lambda)\left(\int_{0}^{\infty} B_{\lambda}(\lambda, T) P(a, T) \mathrm{d} T\right) n(a) \mathrm{d} a \tag{1.37}
\end{equation*}
$$

We now have all the necessary tools to interpret the observations and access the nature of the dust grains present in the ISM.

### 1.2.5 The nature of dust grains unveiled

Now that we know how the physical processes responsible for dust extinction and emission depend on the nature of the dust grain, we can access the chemical composition and size distribution of the dust in the ISM from a well-chosen set of observational constraints.

## Chemical composition

Dust grains are built from the available elements in the ISM, and those are then depleted from the gas phase. The difference between what is expected to be ejected from stars, i.e., the total element abundances, and what is actually observed in the ISM in the gas phase, is called the elemental depletions. Elemental depletion measurements can be used to estimate which metals, i.e., elements heavier than Helium, are locked up into dust and to what fraction. Figure 1.8 shows the elemental depletions for several elements as a function of their condensation temperature, i.e., the ease at which a given element is incorporated into dust. The $\mathrm{C}, \mathrm{O}, \mathrm{Si}, \mathrm{Mg}$ and Fe elements, are among the main constituents of dust grains and are indeed depleted from the gas phase.


Fig. 1.8. Interstellar elemental depletions for a variety of metals as a function of the condensation temperature (from Savage \& Sembach 1996).

As stated throughout this Section, observations of the ISM (such as UV extinction, IR emission, elemental depletion, etc.) provided several clues to determine what kind of grains are present in the ISM. Several types of dust grains have been identified: silicate grains, carbon grains and polycyclic aromatic hydrocarbons (PAHs). Of course, the composition of interstellar dust has been widely debated but these three families are considered to be the most important dust components.

Silicates: The broad band absorption features at 9.7 and $18 \mu \mathrm{~m}$ observed in the interstellar extinction curve (Fig. 1.4) have been attributed to silicate-based material. They are made from SiO blocks and present a wide variety of composition: $\mathrm{MgSiO}_{3},\left(\mathrm{Mg}, \mathrm{Fe}^{2}\right) \mathrm{SiO}_{4}, \mathrm{Mg}_{2} \mathrm{SiO}_{4},(\mathrm{Mg}, \mathrm{Fe})_{2} \mathrm{SiO}_{4}$, etc. The absence of a sub-feature in the $9.7 \mu \mathrm{~m}$, contrary to what laboratory measurements of
absorption profiles show for crystalline silicates, may be the sign that these grains are mostly amorphous in the ISM (except in circumstellar environments) (Kemper et al. 2004). The features at 9.7 and $18 \mu \mathrm{~m}$ are usually attributed to the Si-O stretching and O-Si-O bending modes respectively.

Carbons: Small graphitic grains are required to explain the $2175 \AA$ feature of the observed interstellar extinction curve. Indeed graphite has a strong resonance at $\sim 2000 \AA$ due to $\pi \rightarrow \pi^{\star}$ transition associated with its aromatic bond. The position and profile of this band is determined by several parameters such as grain size, grain shape and grain optical properties. Graphite can evolve into amorphous carbon grains if the physical conditions of their environment are suitable.

Polycyclic Aromatic Hydrocarbons (PAHs): Observations in the 1970s and 1980s have shown that the MIR spectra of bright sources with associated dust and gas are dominated by broad emission features, that could not be explained by the silicate or carbon grains. Spectroscopic studies with ISO have revealed that these broad emission features were dominating the NIR spectrum of a wide variety of objects: from planetary nebulae to (nearby) galaxies (with the exception of asymptotic giant branch stars and deeply embedded young stellar objects). These broad emission features are observed primarily at $3.3,6.2,7.7,8.6$ and $11.3 \mu \mathrm{~m}$ as well as other wavelengths, and were long called the "unidentified infrared bands" (UIB).

Duley \& Williams (1981) showed that the 3.3 and $11.3 \mu \mathrm{~m}$ features could correspond to vibrational modes of C-H groups in aromatic materials. A few years later, Leger \& Puget (1984) proposed large planar molecules made of aromatic cycles to explain these NIR broad emission features.



PAH molecule

Fig. 1.9. (left) Structure of a benzene molecule with the principal vibration modes indicated. (right) Example of a PAH molecule, here benzo(a)pyrene $\mathrm{C}_{20} \mathrm{H}_{12}$.

An example of a PAH molecule is shown in Fig. 1.9, together with the typical vibration modes of benzene: C-H and C-C stretching modes, C-H in/out of the plane bending modes. PAHs typically possess between 50 and 1000 atoms. The main vibration modes give rise to the NIR emission features. Figure 1.10 shows an example of a PAH spectrum in the Orion Bar and we can identify the different bands: the $3.3 \mu \mathrm{~m}$ feature from C-H stretching mode, the $6.2 \mu \mathrm{~m}$ feature from C-C stretching mode, the $7.7 \mu \mathrm{~m}$ feature which is a combination of C-C stretching and C-H in-the-plane bending modes, the $8.6 \mu \mathrm{~m}$ feature from C-H in-the-plane bending mode and the $11.3 \mu \mathrm{~m}$ feature from C-H out-of-the-plane bending mode. The central wavelength of the observed features can vary depending on the structure of the molecules, and the relative intensity of the bands depends on the ionisation degree of the PAHs in the ISM.


Fig. 1.10. Example of a PAH spectrum from the Orion Bar (from Peeters et al. 2002). The main PAH features are indicated in red along with the vibration mode they originate from. Several ionic fine-structure lines have been identified too.

## Size distribution

Since dust is responsible for extinction from UV to visible wavelengths, and the extinction crosssection depends on the grain radius $a$, dust grains must span a range of sizes. Mathis et al. (1977) was the first to propose a size distribution for spherical silicate and carbon grains (hereafter called "MRN" model) that was able to reproduce the interstellar extinction curve in the Galaxy between 0.11 and $1 \mu \mathrm{~m}$. This size distribution was a simple power-law distribution for grain sizes between 5 and 250 nm :

$$
\begin{equation*}
n(a) \propto a^{-3.5} \tag{1.38}
\end{equation*}
$$

where $n(a)$ is the size distribution and $n(a) d a$ is the number of particles with radius between [a, a+da]. The MRN model did not include PAHs and indeed, IRAS observations at 12 and 25 $\mu \mathrm{m}$ revealed that the MRN model could not reproduce the observations of the IR emission from the diffuse Galactic medium. The first model that was able to reproduce both the extinction and emission of the diffuse Galactic ISM was the Desert et al. (1990) model that took into account PAHs, with sizes between 0.4 and 0.12 nm , on top of the silicate and carbon grain populations. To explain the observations, the Desert et al. (1990) model requires dust grains spanning a wide range in grain sizes ( $\sim 3$ orders of magnitude) from sub-nanometre grains to grains of a fraction of a micron. The PAHs are the smallest grains, with sub-nanometre to nanometre sizes. These grains are transiently heated by the absorption of a single photon (see Section 1.2.4) and cool down by emitting MIR photons in vibrational modes representative of their aromatic structure, producing the broad emission features described previously. Another component, very small grains (VSG), is also required to explain the continuum emission in the MIR in the Desert et al. (1990) model. These nanometre grains also undergo significant temperature fluctuations in the ISM. They can be made from either carbonaceous or silicate material. The third component of the dust model is the big grain (BG) component, with sizes $\geq 10 \mathrm{~nm}$ to $\sim 400 \mathrm{~nm}$, necessary to account for the FIR emission. The BG component usually dominates the total dust mass, and is mostly composed of
silicate grains. Depending on their sizes, these bigger grains can be either in thermal equilibrium with the stellar radiation field or stochastically heated by the ISRF.

The grain size distribution we will consider for our study is that presented by Zubko et al. (2004), who simultaneously fit the extinction, emission and also elemental abundances of the diffuse Galactic ISM. The difference with the Desert et al. (1990) model is that Zubko et al. (2004) use both PAHs and graphite grains, and not only graphite, to explain the $2175 \AA$ feature and the FUV continuum of the interstellar extinction curve. We refer to Chapter 7 for a complete description of this grain size distribution and its determination.

### 1.2.6 Life and death of a dust grain in the ISM

As we mentioned in Section 1.2.1, dust is formed from the metals produced by the stars. The main stardust factories are circumstellar shells around evolved stars that provide the necessary high densities and temperatures to produce the dust grains. Two main categories in these dust sources can be distinguished: low-mass stars that undergo quiescent mass-loss and return dust to the ISM via stellar winds; and high-mass stars with violent death that return their ejecta to the ISM via supernova explosions:

- Mass-loss through stellar winds: These stellar sources can be low-mass stars ( $\mathrm{M} \lesssim 8 M_{\odot}$ ) in the red giant phase of their evolution (e.g., M giants, carbon stars, radio luminous $\mathrm{OH} / \mathrm{IR}$ stars), but dust also forms in the circumstellar shells around supergiants, planetary nebulae and Wolf-Rayet stars. Those sources are either C-rich or O-rich. Depending on their C/O abundance ratio, the type of dust produced will differ. For C-rich stars ( $\mathrm{C} / \mathrm{O}>1$ ), all of the oxygen is tied up in CO, and no oxygen is available for the dust. Thus the dust grains are carbon rich ( C and SiC types). For O-rich stars ( $\mathrm{C} / \mathrm{O}<1$ ), the remaining oxygen can combine with other elements, and especially Si to form silicate type material (based on the SiO block). Dust grains are supposed to survive the injection phase in the ISM.
- Explosive mass-loss: $\left(\mathrm{M} \geq 8 M_{\odot}\right)$ Novae, Type Ia and Type II supernovae (SN) are sites of dust production, Type II SN being the most important one. However, most of the dust grains are thought to be immediately destroyed by SN shocks when injected into the ISM.

The main dust production sites are summarised in Table 1.2.
Table 1.2. Summary of the dust production sites and their contribution in the ISM, for the Galaxy (adapted from Jones et al. 1997).

| Source | Type of dust | Contribution $\left[10^{-6} M_{\odot} \mathrm{kpc}^{-2} \mathrm{yr}^{-1}\right]$ |
| :--- | :---: | :---: |
| M giants | Silicate | 3 |
| Radio Loud OH/IR stars | Silicate | 2 |
| Carbon stars | SiC - Carbon - PAHs | 2 |
|  |  |  |
| Supergiants | Silicate | 0.2 |
| Planetary Nebulae | Carbon - PAHs | 0.03 |
| Wolf Rayet stars | Carbon - PAHs | 0.03 |
| Type II SN | Silicate - Carbon | $0.15-14$ |
| Type Ia SN | Silicate - Carbon | $0.03-2.3$ |
| Novae | Silicate - SiC - Carbon - PAHs | $0.003-0.2$ |

After formation, dust is submitted to various evolution processes either constructive (i.e., accretion - coagulation) or destructive processes (such as sputtering, erosion, vaporisation and shattering). These evolution processes result from supernovae shocks that can induce grain-grain collisions or impacts with atoms from the hot gas. This various processes can thus affect the grain size distribution.

Collisions with atoms from the hot gas provoke thermal and chemical sputtering and erosion of the dust grain, transferring grain material back to the gas phase.

Grain-grain collisions have different consequences depending on the velocity of the shock. For high-velocity shocks ( $\mathrm{v} \geq 20 \mathrm{~km} . \mathrm{s}^{-1}$, Tielens et al. 1994), the grains are vaporised and grain material is returned to the gas phase. The destruction of dust grains by SN shocks is the dominant process of recycling dust into gas in the ISM. Intermediate velocity shocks ( $\mathrm{v} \sim 2 \mathrm{~km} . \mathrm{s}^{-1}$, Jones et al. 1996) lead to shattering of the dust grains and transfer mass from big grains to small grains. At low velocities, (v $\lesssim 0.1-1 \mathrm{~km} . \mathrm{s}^{-1}$, Jones 2004) grain-grain collisions lead to grain sticking and coagulation. This impacts the grain size distribution in the opposite way, by increasing the mean grain size. Coagulation mainly occurs in dense regions such as molecular clouds and circumstellar shells.

The typical timescale for dust formation is $\sim 3 \times 10^{9} \mathrm{yr}$ (Jones \& Tielens 1994) whereas the typical timescale for dust destruction is $\sim 5 \times 10^{8} \mathrm{yr}$ (Jones et al. 1996). Because we still observe dust in the ISM, we need re-accretion processes in the ISM to maintain equilibrium between formation and destruction of the dust grains. Dwek (1998) included re-accretion in the ISM in their model for dust evolution in the Galaxy and indeed found that dust formation and re-accretion can be in equilibrium. How dust is preserved for re-accretion is still a debated question.

### 1.2.7 Dust: a fundamental component of the ISM

Even though dust represents only $\sim 1 \%$ in mass in the ISM, it plays a major role in the total emission of a galaxy: in our Galaxy, dust reprocesses about $30 \%$ of the stellar power, and it can be as high as $\sim 99 \%$ in a starburst galaxy (i.e., a galaxy undergoing a burst of star formation). Through the absorption of stellar light, dust is the main source of opacity in the ISM for non-ionising photons.

Dust also impacts the chemical evolution of the ISM by locking up a significant fraction of the heavy elements available in the ISM.

Figure 1.11 illustrates the role of interstellar dust in the lifecycle of the ISM. Stars are born in molecular clouds, form elements in their cores, and re-inject them into the ISM as they evolve and die. Low-mass stars inject elements into the ISM through stellar winds and massive stars through stellar winds and supernovae explosions (Section 1.2.6). Dust grains form from the available elements in the ISM and thus regulate the abundance of elements in the different ISM phases: metals are depleted from the gas phase, when used to form a dust grain. As we saw in Section 1.2.6, through constructive processes (accretion or coagulation) more metals will be locked up in the solid phase of the ISM, while through destructive processes (erosion and sputtering), metals are released once again into the gas phase and are available for the next generation of stars.

As we will see in Section 1.3, dust grains participate actively in the heating and cooling of the ISM: they cool the ISM by absorbing stellar photons and heat the ISM by re-emitting the absorbed power. The main heating process in the neutral ISM is indeed the photoelectric effect on dust grains. FUV photons are absorbed by the dust grain and this generates a diffusion of electrons in the grain. These electrons loose energy through collisions while they diffuse in the grain and eventually escape into the gas phase, thus heating it. The major actors of the photoelectric effect are the smallest grains or very large molecules (PAHs).

Dust also has a double effect on star formation. Indeed to form stars we need cold molecular clouds and thus enough shielding from interstellar radiation to be able to form the molecules and cool the gas. Dust absorbs the stellar radiation that would otherwise dissociate the $\mathrm{H}_{2}$ molecules, and


Fig. 1.11. Schematic view of the processes of matter recycling in the ISM. Credits: F. Galliano.
thus participates actively in the cooling of the ISM. It is also a catalyser of $\mathrm{H}_{2}$ molecule formation by providing a surface where two H atoms can accrete, meet and react. Indeed the presence of dust grains can increase the $\mathrm{H}_{2}$ formation rate by about two orders of magnitudes compared to $\mathrm{H}_{2}$ formation without dust (Tielens 2005), thus activating star formation. Hence dust is an important agent of galaxy evolution.

### 1.3 Gas phases of the ISM

The gas is the other major component of the ISM and is the main fuel for star formation. It represents $99 \%$ of the ISM mass, the majority of which is in the form of hydrogen gas ( $\sim 71.5 \%$ ). The rest of the gas reservoir consists of Helium ( $\sim 27 \%$ ) and gaseous metals ( $\sim 1.5 \%$ ) (values from Asplund et al. 2009). It is convenient to describe the different gas phases of the ISM by the physical state of hydrogen: ionised in the form of $\mathrm{H}^{+}$, neutral in the form of H atoms, molecular in the form of $\mathrm{H}_{2}$. For the different phases, we follow the description of Tielens (2005).

### 1.3.1 Ionised phase

The ionised phase in the ISM is composed of three different regions: the Hot Ionised Medium (HIM), Hir regions and the Warm Ionised Medium (WIM).

## The Hot Ionised Medium - hot and diffuse

Typical temperature $\sim 10^{5}-10^{6} \mathrm{~K}$ - Typical density $\sim 3 \times 10^{-3} \mathrm{~cm}^{-3}$
The HIM consists of very hot coronal gas originating from SN ejected material. It is a very hot but diffuse component of the ionised gas phase of the ISM. This gas is heated and ionised by shock waves generated from supernovae explosions and stellar winds of dying stars. It cools down via thermal X-ray continuum emission and line emission from highly ionised species, such as Civ, Siv, Nv, Ovi or even Ovir and Oviii. In such hot plasmas, several radiation processes can occur: Bremsstrahlung continuum emission (also called "free-free", transition of a free electron between two states), discrete line emission (transition between two levels of the ion), radiative recombination
continuum emission (capture of an electron into a bound state), dielectronic recombination line emission (capture of a free electron into a doubly excited ion state) or two-photon continuum emission (simultaneous emission of two photons) (see Ehle 2005 for a review).


Fig. 1.12. (left) Example of HIM: Supernovae remnants Simeis 147 in the Taurus constellation. The image is a colour composite of 66 colour band images from the Palomar Observatory. (right) Example of an Hir region: the Orion Nebula M42 in the Orion constellation. The image is a colour composite of hydrogen, oxygen and sulfur gas emission (source: APOD).

## HiI regions - hot and dense

Typical temperature $\sim 10^{4} \mathrm{~K}$ - Typical density $\sim 1-10^{5} \mathrm{~cm}^{-3}$
Hir regions are regions around massive O and B stars or OB associations and are thus directly linked to star formation. These stars ionise their immediate surrounding medium creating an ionisation front in the ISM which delineates the Hir region. They span a wide range of densities: from $1 \mathrm{~cm}^{-3}$ for the most diffuse nebulae to $10^{3}-10^{4} \mathrm{~cm}^{-3}$ for compact Hir regions and up to $10^{5}$ $\mathrm{cm}^{-3}$ for the densest regions. Their temperature is directly linked to the gas density but also to the amount of metals in the region, and they are thus extensively used to determine the quantity of metals in the ISM of a galaxy (see Section 2.1, Pilyugin \& Thuan 2005; Izotov et al. 2006). The main heating sources of these regions are photoionising photons (i.e., photons with $h \nu \geq 13.6 \mathrm{eV}$ ) from the OB stars, and dust grains through photoelectric heating. Hir regions cool down over the whole electromagnetic spectrum from UV to radio wavelengths. They are strong sources of thermal radio emission (free-free emission), $\mathrm{H} \alpha$ recombination line emission in the visible or forbidden finestructure line emission in the FIR (such as Oiir). The dust present in these regions also absorbs and scatters stellar light (reflection nebulae) and sometimes can completely obscure the Hir region at visible wavelengths. This absorption results in bright MIR continuum emission observed in these regions arising from the heated dust grains.

Figure 1.12 shows an example of the hot ionised phase as viewed from optical emission lines.

## The Warm Ionised Medium - warm and diffuse

## Typical temperature $\sim 8000 \mathrm{~K}$ - Typical density $\sim 0.1 \mathrm{~cm}^{-3}$

The WIM represents $90 \%$ of the ionised gas phase and is a wide-spread, diffuse medium across galaxies. It is heated by ionising photons that escape HiI regions and/or by UV radiation from isolated hot stars. The cooling processes are the same as in an HII region but the dominant emission
used to probe the WIM is free-free radio emission or $\mathrm{H} \alpha$ recombination line emission. The spectra of diffuse Hir regions differ from those of dense Hir regions by showing strong emission from low stages of ionisation (NiI, Sii) and weak emission from higher stages of ionisation (OiiI). In a dense Hir region, on the other hand, we would rather see the opposite.

### 1.3.2 Atomic phase

The atomic phase is defined as the phase of the ISM where photons have energies lower than 13.6 eV (i.e., non-ionising photons) and where molecules do not exist. In this phase we thus find neutral atomic species (Hi, Ci, Oi) as well as ionised species with ionisation potentials lower than that of hydrogen (i.e., Cir, Sir, Siir). It is composed of the Warm Neutral Medium (WNM) and the Cold Neutral Medium (CNM): the atomic phase can be seen as an ensemble of cold and dense clouds (CNM) bathing in a more diffuse and warmer intercloud phase (WNM).
WNM: Typical temperature $\sim 8000 \mathrm{~K}$ - Typical density $\sim 0.5 \mathrm{~cm}^{-3}$
CNM: Typical temperature $\sim 80 \mathrm{~K}$ - Typical density $\sim 50 \mathrm{~cm}^{-3}$
The neutral gas is usually mapped through the Hı line at 21 cm , arising from the transition between two hyperfine structure levels of the ground state of atomic hydrogen.

The major sources of heating for the WNM are photoelectric heating by dust grains, FUV pumping of $\mathrm{H}_{2}$ (see below), cosmic rays and X-rays. The WNM mainly cools through the optical $\mathrm{H} \alpha$ recombination line.

## Photodissociation Regions (PDRs)

We can define a transition phase that separates the ionised medium from the molecular cloud cores, the photodissociation regions (PDRs), where FUV photons with $6 \mathrm{eV}<\mathrm{h} \nu<13.6 \mathrm{eV}$ control the physics and chemistry of the gas. PDRs contain part of the atomic (the dense and cold neutral medium) and part of the molecular phases of the gaseous ISM, and are directly associated with star formation (to be distinguished from diffuse neutral gas). Dust is associated with these regions and plays an important role in heating up the gas and absorbing the incoming stellar radiation.

Fig. 1.13 shows a schematic representation of a PDR. FUV photons penetrate the cloud and ionise the gas creating an ionisation front at the outer edge of the PDR. Then photons with energies $<13.6 \mathrm{eV}$ penetrate deeper into the cloud ionising other species such as carbon and oxygen, and photodissociating the molecules. This creates an Hi/Cir region. Once the flux of dissociating $\mathrm{H}_{2}$ photons is sufficiently attenuated (usually around $\mathrm{A}_{V} \sim 2 \mathrm{mag}$ ), $\mathrm{H}_{2}$ molecules form in a layer of $\mathrm{H}_{2} /$ CiI. Beyond this layer ( $\mathrm{A}_{V} \sim 4 \mathrm{mag}$ ), carbon-ionising photons are also attenuated and there is a quick transition from Cir to Ci and then to CO. The transition from atomic to molecular oxygen marks the inner edge of the PDR: where the molecular core is shielded from photodissociating photons.

Two major processes dominate the heating of the gas in PDRs: photoelectric heating by the dust grains and FUV pumping of $\mathrm{H}_{2}$. As explained in Section 1.2.7, for the photoelectric effect, an electron is ejected from the dust grain and then heats the gas through collisions. This process is less efficient in PDRs than in Hir regions, as the incoming photons are less energetic. The photoelectric heating efficiency also depends on the grain size and charge. The major actors of photoelectric heating are thus the smallest grains of the ISM, which are often the most numerous, and provide the largest cumulative surface: PAHs and small graphitic grains, emitting in the MIR range. The charge of the dust grain is controlled by the balance between photoionisation and recombination with an electron. This depends on the intensity of the incident radiation field, the temperature of the gas and the density of electrons (through the ionisation parameter $\gamma=G_{0} T_{e}^{0.5} / n_{e}$ where $G_{0}$ is the intensity of the radiation field expressed in Habing units ${ }^{5}$, and $n_{e}$ is the electron density in the

[^4]

Fig. 1.13. Schematic view of a PDR.

PDR, usually $\sim 10^{-4} \mathrm{~cm}^{-3}$ ).
FUV photopumping of $\mathrm{H}_{2}$ occurs through photons with energies $>11.2$ and 12.3 eV (LymanWerner electronic transitions). Absorption of these FUV photons brings the $\mathrm{H}_{2}$ molecule to a bound excited electronic state. The molecule then cascades back to ground vibrational states releasing energy into the gas. In 10 to $15 \%$ of the cases it also leads to $\mathrm{H}_{2}$ dissociation. Other heating sources in PDRs are ionisation of carbon (releasing an electron into the gas), formation of $\mathrm{H}_{2}$ and $\mathrm{H}_{2}$ photodissociation. In denser regions, the contribution of grain-grain collisions to the heating of the gas may become important.

Cooling of the PDR occurs mostly through FIR fine structure line emission of abundant ionic and atomic species i.e., [CiI] $\lambda 158 \mu \mathrm{~m}$ or [OI] $\lambda 63,146 \mu \mathrm{~m}$. PDRs are thus bright in the IR: MIRFIR continuum emission, MIR PAH emission features and FIR fine structure line emission. PDRs are also sources of rotational lines from molecular species, mostly CO , but also $\mathrm{CO}^{+}, \mathrm{CN}, \mathrm{C}_{2}$, and rotational-vibrational transitions of warm $\mathrm{H}_{2}$.

### 1.3.3 Molecular phase

Typical temperature $\sim 10 \mathrm{~K}$ - Typical density $\geq 200 \mathrm{~cm}^{-3}$
The molecular phase is the densest and coldest phase of the gaseous ISM and is defined by every locations where there is molecular hydrogen, which is the most important and abundant molecule in the ISM. Other molecules can also be found deeper into the molecular cloud such as CO, CS, $\mathrm{HCN}, \mathrm{O}_{2}, \mathrm{CH}+, \mathrm{OH}, \mathrm{H}_{2} \mathrm{O}$, etc, if there is enough shielding from stellar radiation. Molecules can survive in these dense clouds because radiation is weakly penetrating the molecular cores. Today, about 200 molecules have been identified in the ISM (Tielens 2005). The molecular phase is in the form of discrete clumps of sizes of the order of tens of parsecs and masses of $10^{3}-10^{6} M_{\odot}$, that are gravitationally-bound rather than in pressure equilibrium with the other phases. They are considered to be the stellar nurseries where stars form by fragmentation processes and gravitational
collapse of the clouds. Figure 1.14 shows a well-known example of a molecular cloud: the Horsehead nebula, from the Orion giant molecular cloud.


Fig. 1.14. Example of a molecular cloud: the famous HorseHead Nebula illuminated from behind by IC434 (source APOD).
$\mathrm{H}_{2}$ molecules, as well as other molecules, are most easily formed on the surface of dust grains at cold temperatures. They can be destroyed by UV photons but are efficient at self-shielding. Molecular gas is mainly heated by cosmic rays that are able to penetrate the clouds. Molecules are excited by collisions or shocks. Molecular clouds mainly cool through vibrational and rotational molecular transitions, CO being the dominant coolant. Dust is cold in molecular clouds and mostly emits in the FIR-to-submm wavelengths.
$\mathrm{H}_{2}$ is a symmetric molecule and does not have dipolar momentum, rendering it very difficult to observe directly through its rotational lines. Some electronic transitions can be observed along the line of sight towards hot stars. NIR rotational vibrational transitions of $\mathrm{H}_{2}$ are indicative of warm ( $\sim 1000 \mathrm{~K}$ ) and dense ( $\sim 10^{4} \mathrm{~cm}^{-3}$ ) gas in shocks. This warm $\mathrm{H}_{2}$ represents a small mass fraction of the total molecular gas mass. Cold $\mathrm{H}_{2}$ gas is indirectly observed, through CO rotational transitions in the submm: $\mathrm{CO}(1-0)$ at $2.6 \mathrm{~mm}, \mathrm{CO}(2-1)$ at $1.3 \mathrm{~mm}, \mathrm{CO}(3-2)$ at $870 \mu \mathrm{~m}$, etc., as CO is the most abundant molecule after $\mathrm{H}_{2}$.

Converting the observed CO intensity into a molecular gas mass is usually done via a conversion factor, $X_{C O}$, defined by:

$$
\begin{equation*}
X_{C O}=N\left(H_{2}\right) / I_{C O} \tag{1.39}
\end{equation*}
$$

where $N\left(H_{2}\right)$ is the column density of $\mathrm{H}_{2}$ we want to determine, and $I_{C O}$ is the CO intensity measured from one given transition. $X_{C O}$ was first determined from $\mathrm{CO}(1-0)$ observations in the Galaxy, calibrated from cold clouds assumed to be virialised. The Galactic conversion factor, $X_{C O, M W}$, has been estimated to be $2.0 \times 10^{20} \mathrm{~cm}^{-2}\left(\mathrm{~K} \mathrm{~km} \mathrm{~s}^{-1}\right)^{-1}$ (Ackermann et al. 2011). Sometimes the conversion factor is written in terms of mass surface density: $\alpha_{C O}$, with:

$$
\begin{equation*}
\alpha_{C O}=M\left(H_{2}\right) / L_{C O} \tag{1.40}
\end{equation*}
$$

where $L_{C O}$ is the CO luminosity and $M\left(H_{2}\right)$ is the molecular gas mass. $\alpha_{C O}$ is related to $X_{C O}$ via:

$$
\begin{equation*}
\alpha_{C O}=X_{C O} \times m_{H_{2}} \times \frac{\Omega}{4 \pi D^{2}} \tag{1.41}
\end{equation*}
$$

where $m_{H_{2}}$ is the mass of one $\mathrm{H}_{2}$ molecule in $M_{\odot}, \Omega$ is the source solid angle and $D$ is the distance to the source. $\alpha_{C O}$ is expressed in units of $M_{\odot}\left(\mathrm{K} \mathrm{km} \mathrm{s}^{-1} \mathrm{pc}^{2}\right)^{-1}$, and the Galatic value is $3.8 M_{\odot}\left(\mathrm{K} \mathrm{km} \mathrm{s}^{-1} \mathrm{pc}^{2}\right)^{-1}$.

This conversion between CO intensity and molecular gas mass is still under investigation as $X_{C O}$ strongly depends on various parameters such as the radiation field intensity, metallicity, gas density and velocities, etc. (Shetty et al. 2011; Feldmann et al. 2012). The impact of metallicity on the $X_{C O}$ conversion factor is discussed in Chapter 2. For a review on $X_{C O}$, see Bolatto et al. (2013).

Table 1.3 summarises the main characteristics of the various gas phases in the Galaxy.
Table 1.3. Characteristics of the phases of the ISM in the Galaxy (adapted from Tielens 2005). For each phase, a typical temperature and density are given, along with its total mass in the Galaxy and the local surface density in the solar neighbourhood. Means of detection of each phase are given in the last column.

| Phase | $\mathrm{T}^{a}[\mathrm{~K}]$ | $\mathrm{n}^{b}\left[\mathrm{~cm}^{-3}\right]$ | $\mathrm{M}^{c}\left[10^{9} M_{\odot}\right]$ | $\Sigma^{d}\left[M_{\odot} \mathrm{pc}^{-2}\right]$ | Observation |
| :--- | :---: | :---: | :---: | :---: | :--- |
| Hot Ionised Medium | $10^{5}-10^{6}$ | 0.003 | - | 0.3 | Thermal X-ray emission <br> Ionised metals absorption/emission lines <br> Thermal radio emission |
| HiI regions | $10^{4}$ | $1-10^{5}$ | 0.05 | 0.05 | Thermal radio emission <br> $\mathrm{H} \alpha$ recombination line <br> FIR fine-structure lines <br> Bright MIR continuum emission from dust |
| Warm Ionised Medium | 8000 | 0.1 | 1.0 | 1.1 | H $\alpha$ recombination line <br> Thermal radio emission |
| Warm Neutral Medium | 8000 | 0.5 | 2.8 | 1.5 | HI 21 cm absorption/emission line |

${ }^{a}$ Typical temperature.
${ }^{b}$ Typical density.
${ }^{c}$ Representative mass in the Milky Way.
${ }^{d}$ Mass surface density in the solar neighbourhood.

### 1.4 Probing the ISM of galaxies

### 1.4.1 Disentangling the different components

Sections 1.2 and 1.3 give us means to interpret the SED of a galaxy and we can now disentangle the various physical components and processes at the origin of what is observed in galaxies: i.e., decompose the blue profile in Fig. 1.2.

Fig. 1.15 shows the detailed decomposition of the total SED of a galaxy in the different contributions from stars, dust and gas ${ }^{6}$. The observational constraints are indicated in black and are broad-band photometry from space or ground-based telescopes, spectrum (in the MIR) and emission lines. The total modelled emission of the galaxy that best reproduces the observational constraints is shown with the dashed grey curve. This total emission can be decomposed into the contributions of the various physical components in the galaxy. The most massive stars (in cyan) ionise the gas and heat the dust surrounding them. The emission from this gas is shown in yellow: note all of the emission lines, due to the ionised species in the gas. The emission from this dust is shown in blue: note the peak emission in the MIR, around $30 \mu \mathrm{~m}$, indicative of a hot dust component with high equilibrium temperatures and the absence of PAH features in this particular case (PAHs have been photodissociated by the hard and intense FUV photons in these regions). These combined emission profiles correspond primarily to the emission from Hir regions in the galaxy.


Fig. 1.15. Detailed view of the SED of a galaxy with the contributions from stars (in cyan and red), dust (in pink, green and blue) and gas (in yellow and brown). The total SED is the dashed grey line. Observations are the black points and spectrum.

Emission from older and less massive stars is shown in red, and dust heated by this radiation is shown in pink: note that here we do see the PAH emission features in the MIR. The peak of the dust emission is shifted to longer wavelengths, around $90 \mu \mathrm{~m}$, indicative of colder grains. This corresponds to the emission from the neutral phase of the ISM. The emission from even colder dust embedded in molecular clouds, far away from any heating source, is shown in green with a peak in the submm domain ( $\sim 300 \mu \mathrm{~m}$ ). Finally the radio emission is due to emission from accelerated

[^5]charges in the ISM (free-free and synchrotron radiation), originating in various regions, namely $\mathrm{H}_{\mathrm{II}}$ regions and the hot gas (HIM). Absorption of stellar light by dust is shown in the UV-visible by the stripped areas in the stellar emission. Note also the Lyman edge at $912 \AA$ where the total emitted luminosity of the galaxy drops by about two orders of magnitude.

A wide range of instruments is available to probe the numerous components of the ISM of galaxies over all wavelengths, thus allowing us to obtain the necessary observational constraints. Here we present some of the telescopes and instruments that will be used in this thesis.

GALEX (GALaxy evolution EXplorer) is a space UV telescope observing at 1539 and $2316 \AA$.
2MASS (Two Microns All Sky Survey) is a large ground-based survey in the J ( $1.25 \mu \mathrm{~m}) \mathrm{H}(1.65$ $\mu \mathrm{m})$ and $\mathrm{K}(2.20 \mu \mathrm{~m})$ bands. It was conducted from the Mt. Hopkins Telescope in Arizona (USA) for the Northern hemisphere and from the Cerro Tololo Interamerican Observatory (in Chile) for the Southern hemisphere.

ISO (Infrared Space Observatory) contains an infrared camera (ISOCAM) observing between 2.5 and $17 \mu \mathrm{~m}$.

Spitzer is a space IR telescope containing three instruments: IRAC and MIPS, both photometers observing from 3.6 to $160 \mu \mathrm{~m}$, and IRS, the spectrometer covering 7 to $40 \mu \mathrm{~m}$. A more complete description of these instruments is given in Chapter 6.

Herschel carries three instruments, PACS, SPIRE and HIFI, observing from 55 to $670 \mu \mathrm{~m}$ in FIR and submm wavelengths. PACS and SPIRE are both photometers and medium-resolution spectrometers and HIFI is a high-resolution spectrometer. Herschel is at the very heart of our study and the second part of this manuscript (Chapters 3 to 5) is entirely dedicated to Herschel data.

JCMT, APEX, IRAM (James Cerk Maxwell Telescope - Atacama Pathfinder EXperiment - Institute for Radio Astronomy at Millimetre wavelengths) are three ground-based facilities observing in the submm wavelength range with the instruments SCUBA (Submillimetre Common-User Bolometer Array, JCMT) in Mauna Kea (Hawaii, USA), LABOCA (Large Apex BOlometer CAmera, APEX) in Chajnantor (Chile) and MAMBO(Max-Planck-Millimetre-Bolometer, IRAM) in Grenada (Spain). SCUBA observed at 450 and $850 \mu \mathrm{~m}$ and is now replaced by SCUBA2, LABOCA observes at $870 \mu \mathrm{~m}$ and MAMBO observed at 1.2 mm . These three instruments are described in more details in Chapter 6.

Planck is another space telescope with two instruments, LFI and HFI (Low and High Frequency Instruments), observing submm to cm wavelengths from $350 \mu \mathrm{~m}$ to 1 cm .

ALMA (Atacama Large Millimetre/submillimiter Array) is a ground-based radio interferometer observing from $315 \mu \mathrm{~m}$ to 7.5 mm .

### 1.4.2 A wide range of SED shapes

Galaxies present a wide variety of morphological types and can be divided, in simplified terms, into three main groups: spiral galaxies, elliptical and lenticular galaxies, and irregular galaxies. Observations of different types of galaxies have shown that their SED shapes present different characteristics depending on several parameters such as their sizes, luminosities, masses, morphological
types, stellar activities, metallicities, etc., resulting in different relative proportions of the various ISM components. Our Galaxy is a classical barred spiral galaxy (SBc). Schematically spiral and irregular galaxies are more gas-rich, have a relatively young stellar population and can undergo bursts of star formation. On the contrary elliptical galaxies are neutral gas and dust deficient and their stellar population is quite old.

Figure 1.16 shows SEDs of five galaxies of different morphological types and stellar activities. We saw in Sections 1.2 and 1.3 how the SED depends on the ISM properties and Fig. 1.15 gives us the relative contribution of each ISM component. Let us now see on Fig. 1.16 what we can learn for different types of galaxies from the comparison of the different observed SED shapes, using the spiral type as a reference type.


Fig. 1.16. SEDs for various galaxy types (from Madden et al. 2013): Spiral (NGC6946, in black), Dwarf (NGC1705, in purple), Dwarf Elliptical (NGC205, in green), Starburst (M82, in blue) and ULIRG (Arp220, in red). The differences in their SED shapes are due to the variations in the relative contribution of each ISM phase to the total emission of the galaxy. The grey stripes indicate the wavelength coverage of the 6 Herschel bands (see Chapter 3).

Spiral galaxies These galaxies usually contain large quantities of gas and dust and thus show strong dust continuum emission in the IR as well as strong emission lines. Spirals are composed of a rotating disk of stars and ISM (gas and dust), a bulge and sometimes a bar. Usually the bulge shows an older stellar population than the disk. Numerous star-forming regions can be found in the spiral arms. We see on Fig. 1.16, for the spiral NGC6946, the strong PAH features, arising from the active star-forming regions associated with PDRs in the disk, and FIR dust continuum emission
peaking around $100 \mu \mathrm{~m}$.
Elliptical galaxies These galaxies present elliptical light radial profiles. They are known to possess a sparse and dust-poor ISM (Combes 1991) and an old stellar population. Molecular gas is rare, if detected at all, in elliptical galaxies (Sage et al. 2007), and thus they harbour a very low star formation activity. They normally show a weak dust component in their SED compared to the NIR peak, representative of the old stellar component.

Lenticular galaxies They are galaxies possessing a disk and a bulge but no spiral arms. They were originally misclassified by Hubble as intermediate between the roundest elliptical galaxies and spiral galaxies. They are now thought to be former spiral galaxies from which some gas has been removed by interaction with hot cluster gas in galaxy clusters (van den Bergh 2009). Their SEDs indeed show non-negligible dust and PAH emission.

Irregular / Dwarf galaxies As stated by their names, those galaxies are usually very small and do not possess a regular shape, making them difficult to classify. As we see in Fig. 1.16 the dwarfs are much fainter than the "normal" type galaxies, show a FIR peak of the SED shifted to shorter wavelengths, a steeper MIR continuum rise and weaker PAH features. Dwarf galaxies are presented in more details in Chapter 2 as they are the main focus of this work. In Fig. 1.16 we show the SED of NGC1705 which is a dwarf spiral galaxy and NGC205 a dwarf elliptical, and we clearly see the decreased contribution of the dust in the elliptical dwarf compared to the more active star-forming dwarf.

Besides classifying galaxies by morphological type, we can also distinguish galaxies based on their star formation activity:

Starburst galaxies are galaxies that are very actively forming stars (up to $1000 \mathrm{M}_{\odot} \mathrm{yr}^{-1}$ ). To sustain such high star formation rates they consume their gas reservoirs very rapidly and the starburst phase only lasts a few tens of Myr. An important young stellar population creates numerous bright Hir regions surrounded by PDRs. The harder and more intense radiation field they generate leads to a warmer dust component. This translates into a SED with prominent PAH features (from the PDRs) and a dust peak shifted to MIR wavelengths (see M82 in Fig. 1.16).

Active Galactic Nucleus galaxies - AGNs - These galaxies possess a very compact and luminous nucleus which dominates the total luminosity of the galaxy. This enhanced central luminosity is due to the accretion processes occurring in the accretion disk around a central super massive black hole. There are various categories of AGNs: Low-Ionisation Nuclear Emission-line Regions (LINERs galaxies), Seyfert galaxies, quasars, blazars and radio galaxies. Unified models propose that all of these different types are in fact the same object observed from a different point of view (Urry \& Padovani 1995). Among the spectral characteristics of AGN is a "NIR bump" around 2 $10 \mu \mathrm{~m}$ due to very hot dust in the accretion disk heated by the AGN.
(Ultra)-Luminous InfraRed Galaxies - (U-)LIRGs - These galaxies are extremely luminous in the IR wavelength range ( $L_{T I R}>10^{11} L_{\odot}$ for a LIRG and $>10^{12} L_{\odot}$ for a ULIRG) radiating more than $90 \%$ of their energy in the IR. LIRGs are extreme starburst and some also contain an AGN. ULIRGs are thought to be major mergers of gas-rich galaxies and they usually show strong emission lines in their spectra. They may be progenitors of elliptical galaxies (Kormendy \& Sanders 1992; Genzel et al. 2001). Note the prominent MIR-to-FIR peak in the SED of Arp220, a typical ULIRG, in Fig. 1.16. The dust peak is shifted to shorter wavelengths compared to that of spirals

- a sign of hotter dust. The PAH features are absent, because the PAHs can not survive the hard radiation in the vicinity of the AGN.

The properties of irregular / dwarf galaxies, interpretation of their observed SEDs and how they vary compared to "normal" spiral galaxies together with determination of the ISM properties that are responsible for these variations, are fundamental to this thesis work. We thus now focus on a description of these specific environments for the next Chapter.

## Chapter 2

## Low-metallicity galaxies

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In this Chapter, we focus on the characteristics of low-metallicity systems. First we define the metallicity of a galaxy and describe the importance of this parameter in characterising the ISM, before reviewing different ways of determining metallicity in Section 2.1. Physical properties of low-metallicity environments are presented in Section 2.2, where we outline the peculiarities of the ISM in these galaxies and the various motivations for studying them.

### 2.1 Metallicity

### 2.1.1 Definition

Metallicity is defined by the mass fraction of metals in the ISM, where metals encompass all elements heavier than helium. Metallicity is usually noted with the letter " $Z$ ' while " $X$ " and " $Y$ " are the mass fractions of hydrogen and helium respectively. We thus have:

$$
\begin{equation*}
X+Y+Z=1 \tag{2.1}
\end{equation*}
$$

The metallicity is then a measure of the fraction of the initial hydrogen and helium that has been converted into heavier elements. For the lightest elements, up to ${ }^{7} \mathrm{Be}$, this conversion is done during primordial nucleosynthesis, shortly after the Big Bang. Elements up to iron are synthesised in stars, and released in the ISM when stars die (see Section 1.2.6), and elements heavier than iron are produced during supernovae events.

Primordial mass fractions (denoted by " $P$ ", values from Pagel 1997) have evolved to present-day values in the solar neighbourhood (denoted by $\odot$, Asplund et al. 2009) :

$$
\left\{\begin{array} { r l } 
{ X _ { P } } & { \sim 0 . 7 6 }  \tag{2.2}\\
{ Y _ { P } } & { \sim 0 . 2 4 } \\
{ Z _ { P } } & { = 0 }
\end{array} \Rightarrow \left\{\begin{array}{rl}
X_{\odot} & \sim 0.7154 \\
Y_{\odot} & \sim 0.2703 \\
Z_{\odot} & \sim 0.0142
\end{array}\right.\right.
$$

The hydrogen fraction has slowly decreased since the Big Bang as it is converted to heavier elements: helium or metals.

Usually the metallicity is approximated by the ratio of the abundance of one metal element, e.g., oxygen or iron, to the hydrogen abundance, $(\mathrm{O} / \mathrm{H})$ or $(\mathrm{Fe} / \mathrm{H})$, and expressed as $12+\log (\mathrm{O} / \mathrm{H})$ (or $12+\log (\mathrm{Fe} / \mathrm{H})$ ). Note that abundances are in number and not in mass. We use metallicity in terms of oxygen abundance throughout this manuscript. In the solar neighbourhood, we have $(\mathrm{O} / \mathrm{H})_{\odot}=(4.90 \pm 0.56) \times 10^{-4}$ i.e., $12+\log (\mathrm{O} / \mathrm{H})_{\odot}=8.69 \pm 0.05$ (Asplund et al. 2009).

We often talk about galaxies with a given fraction, $f$, of the solar metallicity and note it as " $Z=f Z{ }_{\odot}$ ". It actually means that the oxygen abundance in the galaxy is $f$ times the oxygen abundance of the Galaxy $\left(\mathrm{O} / \mathrm{H}=f(\mathrm{O} / \mathrm{H})_{\odot}\right)$, and often it is assumed that the oxygen abundance scales with the total metal abundance: $(\mathrm{O} / \mathrm{H}) /(\mathrm{O} / \mathrm{H})_{\odot} \sim Z / Z_{\odot}$.

For a galaxy, the metallicity thus traces the history of the stellar activity, i.e., schematically how many cycles of star formation it already went through (see Section 1.2.7 and the lifecycle of matter in the ISM). Metallicity is thus a key parameter of the evolution of galaxies, and is expected to increase with age as the galaxy undergoes chemical enrichment. Investigations of relations between metallicity and other physical properties of galaxies can thus place important constraints on the physical processes governing galaxy evolution. In this work, we especially focus on the relation between metallicity and dust, another key agent of galaxy evolution (Section 1.2.7).

### 2.1.2 Determination

There are several possible ways to determine the metallicity of a galaxy: spectroscopy of ionised gas in Hir regions (Izotov et al. 2006) or planetary nebulas (Pagel 1997); photometry of resolved
stellar populations $((\mathrm{Fe} / \mathrm{H})$ determined by the colour of the red giant branch, e.g., Lee et al. 1993); spectroscopy of individual stars (only possible in the nearest galaxies where individual stars can be resolved, e.g., Haser et al. 1998 in the Magellanic Clouds); spectroscopy or photometry of integrated light for remote galaxies without Hir regions (Worthey 1994); spectroscopy of the cold neutral medium (using absorption lines, Kunth et al. 1994); or X-ray observations of the hot ionised medium (Persic et al. 1998). We refer to Kunth \& Östlin (2000) for a review of the various ways of accessing the metallicity of a galaxy, and detail here only the first: spectroscopic measurements of emission lines in Hir regions.

The abundance of elements in Hir regions depends on the properties of the gas, namely the electron density, $n_{e}$, and the electron temperature, $T_{e}$. These two parameters determine the ionisation stages that can be found in the region for a given element. Several methods exist to derive abundances in Hir regions and we describe here two most commonly used methods: the "direct" method and the "strong-line" method.

## The "direct" method

This method is called "direct" because it allows the direct determination of the electron density and temperature through carefully chosen optical line ratios, after correcting the lines for extinction. We review here the physics allowing us to determine $n_{e}$ and $T_{e}$ from line ratios in the simple case of a two-level system. Let us consider an element collisionally excited to an upper energy level that cascades down to the lower energy level by collisional de-excitation and spontaneously emits a photon. The populations of the two levels can be found by solving the statistical equilibrium equation:

$$
\begin{equation*}
n_{e} n_{l} \gamma_{l u}=n_{e} n_{u} \gamma_{u l}+n_{u} A_{u l} \tag{2.3}
\end{equation*}
$$

where $n_{u}$ and $n_{l}$ are the population densities of the upper and lower energy level respectively; $\gamma_{u l}$ and $\gamma_{l u}$ are the collisional rate coefficients between the two levels; and $A_{u l}$ is the Einstein coefficient for spontaneous emission. In thermal equilibrium, the collisional excitation rate coefficient, $\gamma_{l u}$, is related to the collisional de-excitation rate coefficient, $\gamma_{u l}$, by:

$$
\begin{equation*}
\gamma_{l u}=\gamma_{u l} \frac{g_{u}}{g_{l}} e^{\frac{-E_{u l}}{k T_{e}}} \tag{2.4}
\end{equation*}
$$

where $g_{u}$ and $g_{l}$ are the statistical weights of the two levels; $E_{u l}$ is the energy difference between the two levels; and $k$ is the Boltzmann constant. $\gamma_{u l}$ is given by:

$$
\begin{equation*}
\gamma_{u l}=\left(\frac{2 \pi}{k T_{e}}\right)^{1 / 2} \frac{\hbar^{2}}{m_{e}^{3 / 2}} \frac{\Omega(u, l)}{g_{u}} \tag{2.5}
\end{equation*}
$$

where $m_{e}$ is the mass of the electron; $\Omega(u, l)$ is the collision strength for upward collisions; and $\hbar$ the Planck constant, $h$, divided by $2 \pi$.

With Eqs. 2.3 and 2.4 we can get:

$$
\begin{equation*}
\frac{n_{u}}{n_{l}}=\frac{g_{u}}{g_{l}} e^{\frac{-E_{u l}}{k T_{e}}}\left(\frac{1}{1+\frac{n_{c}}{n_{e}}}\right)=\frac{n_{e} \gamma_{l u}}{A_{u l}}\left(\frac{1}{1+\frac{n_{e}}{n_{c}}}\right) \tag{2.6}
\end{equation*}
$$

where we introduce $n_{c}$, the critical density: $n_{c}=A_{u l} / \gamma_{u l}$. The critical density determines if collisions dominate the de-excitation process. This is the case when $n_{e} \gg n_{c}$. When $n_{e} \ll n_{c}$, collisional de-excitation of the upper level is negligible.
The line intensity that we observe is directly proportional to the cooling rate $\Lambda$, given by:

$$
\begin{equation*}
\Lambda=n_{u} A_{u l} h \nu_{u l} \tag{2.7}
\end{equation*}
$$

with $h \nu_{u l}=E_{u l}$.
In the high-density limit, the observed intensities are independent of the electron density and depend only on the electron temperature. Indeed for $n_{e} \gg n_{c}$, the cooling rate can be expressed as:

$$
\begin{equation*}
\Lambda=n_{l} \frac{g_{u}}{g_{l}} e^{\frac{-E_{u l}}{k T_{e}}} A_{u l} h \nu_{u l} \tag{2.8}
\end{equation*}
$$

On the contrary, in the low-density limit, the observed intensities depend on the electron density. Indeed for $n_{e} \ll n_{c}$, the cooling rate is:

$$
\begin{equation*}
\Lambda=n_{e} n_{l} \gamma_{l u} h \nu_{u l} \tag{2.9}
\end{equation*}
$$

$\Lambda$ depends on $T_{e}{ }^{-1 / 2}$ via the collisional rate coefficient $\gamma_{l u}$ (through Eqs. 2.4 and 2.5). Thus the electron temperature can be determined by using two lines with very high critical densities. The typical line ratio used for the determination of the electron temperature is [OiII] $\lambda 4363 /[\mathrm{OiII}] \lambda \lambda 4959,5007$, because these lines have critical densities of $2.4 \times 10^{7} \mathrm{~cm}^{-3}$ and $7 \times 10^{5} \mathrm{~cm}^{-3}$ respectively.

The electron density is usually measured combining two lines with different critical densities. In between the two critical densities, the line ratio is a good measure of the electron density as one line is in the high-density limit (with intensity $\propto n_{l}$ ) and the other line is in the low-density limit (with intensity $\propto n_{e} \times n_{l}$ ). Line pairs that are commonly used are the [SII] $\lambda 6717 / \lambda 6731$ ratio with respective critical densities of $1.3 \times 10^{3} \mathrm{~cm}^{-3}$ and $3.6 \times 10^{3} \mathrm{~cm}^{-3}$, or [OII] $\lambda 3726 / \lambda 3729$ with $n_{c}$ of $10^{4} \mathrm{~cm}^{-3}$ and $3 \times 10^{3} \mathrm{~cm}^{-3}$.

Once $n_{e}$ and $T_{e}$ are known, the ionic abundance of an element can be derived from:

$$
\begin{equation*}
\frac{N\left(X^{i}\right)}{N\left(H^{+}\right)}=\frac{I_{\lambda}}{I_{H \beta}} \frac{\epsilon_{H \beta}}{\epsilon_{\lambda}} \tag{2.10}
\end{equation*}
$$

where $N\left(X^{i}\right)$ is the ionic abundance of the specie of interest; $I_{\lambda}$ is the measured intensity of the ionic emission line; $I_{H \beta}$ is the measured intensity of the $\mathrm{H} \beta$ recombination line and $\epsilon_{\lambda}$ and $\epsilon_{H \beta}$ are the emissivities of the respective lines. For the ion we are interested in, $\epsilon_{\lambda}$ is directly proportional to the cooling rate and thus to the number density of the ion. The formula to derive the $\mathrm{H} \beta$ emissivity is given in Aller (1984).

The total abundance of the element is obtained by summing the ionic abundances of all of the ionisation stages. As all of the ionisation stages can not always be observed for an element, one has to correct for the unobserved ions via the ionisation correction factors (ICFs) for each element. Fortunately for oxygen, the most abundant ions, $\mathrm{O}^{+}$and $\mathrm{O}^{2+}$, are seen in the optical spectra of $\mathrm{H}_{\text {II }}$ regions and only a small fraction is in the form of $\mathrm{O}^{3+}$. In most cases, the abundance for oxygen can thus be determined directly by:

$$
\begin{equation*}
\frac{\mathrm{O}}{\mathrm{H}}=\frac{\mathrm{O}^{+}}{\mathrm{H}^{+}}+\frac{\mathrm{O}^{2+}}{\mathrm{H}^{+}} \tag{2.11}
\end{equation*}
$$

and this is one of the reasons why the metallicity of a galaxy is usually given in terms of oxygen abundance. To derive ICFs for other elements, we need to know the ionisation structure of the Hir region, that can be derived using photoionisation models.

This "direct" method is the most direct way to access the metallicity of a galaxy but is subject to a number of caveats, among which are:

- The [OIII] $\lambda 4363$ line, necessary to determine $T_{e}$ (OIII), is very weak and most of the time can not be observed easily in high-metallicity environments (e.g., Garnett et al. 2004).
- In high-metallicity regions, temperature fluctuations or gradients in the Hir region may cause a misestimation of the electron temperature. $T_{e}(\mathrm{OIII})$ is then not representative of the true electron temperature in the Hir region leading to a systematic bias in the metallicity determination (usually underestimation, see review by Stasińska 2005; Bresolin 2007).

Other auroral lines detectable at high metallicities (such as [ $\mathrm{NII}^{2}$ ] $\lambda 755$ or $[\mathrm{SII}] \lambda 6312$ ) or high signal-to-noise spectra can overcome the weakness of the [OIII] $\lambda 4363$ line and robust metallicities can be obtained up to roughly solar metallicity.

We refer to Izotov et al. (2006) for an example of determining the metallicity with the "direct" method. Izotov et al. (2006) use the [SII] $\lambda 6717 / \lambda 6731$ ratio to determine the electron density and [OIII] $\lambda 4363 /[\mathrm{OIII}] \lambda \lambda 4959,5007$ to derive the electron temperature. However, the electron temperature is not the same in the low- and high-ionisation zones of the Hir region, especially in Hir regions with low metallicities (Stasińska 1990). Only $T_{e}$ (Oiir) is derived directly from the observations. To get the electron temperature of the different ions of interests, Izotov et al. (2006) use the most recent results from the Hir photoionisation models of Stasińska \& Izotov (2003) to get relations between $T_{e}$ (OiiI) and the temperatures characteristic of other ions. Updated relations from Pagel et al. (1992) are then used in Izotov et al. (2006) to derive ionic abundances from the corresponding electron temperatures, and are based on the three-level-atom solutions of McCall (1984). The ICFs for elements other than oxygen are determined from photoionisation models and are expressed as functions of $\mathrm{O}^{+} /\left(\mathrm{O}^{+}+\mathrm{O}^{2+}\right)$ and $\mathrm{O}^{2+} /\left(\mathrm{O}^{+}+\mathrm{O}^{2+}\right)$.

We will follow the method of Izotov et al. (2006) to determine the metallicity of the galaxies in our study (see Chapter 3). With this method, Izotov et al. (2012) determined the metallicity of 42 extremely low-metallicity galaxies (down to $12+\log (\mathrm{O} / \mathrm{H})=7.12$ ) and compiled a sample of the 17 most metal-poor galaxies known in the local Universe. Their main finding is that there seems to be a metallicity floor around $12+\log (\mathrm{O} / \mathrm{H}) \sim 6.9$, below which no galaxies are found in the local Universe, suggesting that the matter from which dwarf galaxies have formed had been previously pre-enriched to that level by (e.g.,) Population III stars.

## The "strong-line" method

This method is used when the $[\mathrm{OiII}] \lambda 4363$ line is not detected, for example at higher metallicities in low-excitation Hir regions. Empirical metallicity calibrations were derived by fitting the relationship between "direct" metallicities and strong-line ratios for Hir regions. Typical ratios are mostly optical line ratios: [NII] $\lambda 6584 / \mathrm{H} \alpha$ (Pettini \& Pagel 2004), ([OII] $\lambda 3727 / \mathrm{H} \beta$ )/([NII] $\lambda 6584 / \mathrm{H} \alpha$ ) (Pettini \& Pagel 2004), or the " $R_{23}$ " ratio (Pilyugin 2001; Pilyugin \& Thuan 2005; Liang et al. 2007; Yin et al. 2007):

$$
\begin{equation*}
R_{23}=\frac{[\mathrm{O} I I] \lambda 3727+[\mathrm{OIII}] \lambda \lambda 4959,5007}{H \beta} \tag{2.12}
\end{equation*}
$$

The bulk of cooling in dusty environments is done by the FIR fine structure lines while it is done by high-excitation optical lines in low-metallicity regions. Thus we expect $R_{23}$ to increase as the metallicity decreases. Figure 2.1 shows the $R_{23}$ ratio as a function of $12+\log (\mathrm{O} / \mathrm{H})$, and we observe the expected trend for high metallicities. At lower metallicities, [OiII] $\lambda \lambda 4959,5007$ is not dominating the cooling anymore and the relation turns around at $\sim 10 \%$ of the solar metallicity, creating a degeneracy in $R_{23}$. This degeneracy can be lifted using another line ratio, for example [NII] $\lambda 6584 /[\mathrm{OII}] \lambda 3727$. The division between the upper and lower branch of the $R_{23}$ calibration occurs around $\log \left(\left[\mathrm{NII}^{\prime}\right] \lambda 6584 /[\mathrm{OII}] \lambda 3727\right) \sim-1.2$ (Kewley \& Ellison 2008). The "strong-line" method suffers from the same caveats as the "direct" method since the calibration is based on "direct" metallicities.


Fig. 2.1. Oxygen abundance, $12+\log (\mathrm{O} / \mathrm{H})$, as a function of $R_{23}$, derived from McGaugh (1991) models. The figure has been taken from Maier et al. (2004) to illustrate the degeneracy in the $R_{23}$ calibration.

As we are spanning a wide range in metallicity in the sample we use for our study, we will also use the "strong-line" method with the calibration of Pilyugin \& Thuan (2005) to determine the metallicity of our galaxies (see Chapter 3).

Other strong-line $R_{23}$ calibrations can be derived from photoionisation models or from combining models and observations. We refer to Kewley \& Ellison (2008) for a review on the various ways of determining the metallicity of a galaxy through spectroscopic observations of Hir regions.

### 2.2 Dwarf galaxies

### 2.2.1 Definition and classification

## General characteristics

Dwarf galaxies are commonly defined as being small in size with a low mass, a low luminosity and a low metallicity. Where to draw the line between "normal" galaxies and dwarf galaxies is still widely debated but we adopt here the luminosity criterion of Grebel et al. (2003): $M_{V} \geq-18 \mathrm{mag}$, where $M_{V}$ is the absolute magnitude in the V band. A luminosity criterion is often used rather than a mass criterion as luminosity is directly observable for a galaxy, contrary to its mass, and allows a more homogeneous classification. In terms of mass, galaxies with masses smaller than $10^{10} M_{\odot}$ can be considered as dwarf galaxies. Dwarf galaxies in the local Universe can have luminosities as low as $10^{5} L_{\odot}$, masses as low as $10^{7} M_{\odot}$, metallicities as low as $1 / 50 Z_{\odot}$ and sizes in terms of $R_{25}{ }^{1}$ smaller than 5 kpc . For comparison, the Milky Way has $M_{V} \sim-21 \mathrm{mag}$, a total mass, $M_{\text {tot }}$, of $\sim 10^{12} M_{\odot}$ (including dark matter), a total luminosity of $\sim 7 \times 10^{10} L_{\odot}$ (see Table 1.1) and a radius of $\sim 18$ kpc. The stellar content of dwarf galaxies is also low, $\lesssim 10^{9} M_{\odot}$, two orders of magnitude below the stellar mass for the Milky Way. They form stars either via episodic bursts of star formation or continuously (see Section 2.2.2). Dwarf galaxies are the dominant population of galaxies in the Universe (Marzke \& da Costa 1997) even though their low luminosities often restrict their detailed

[^6]study to the nearby Universe. For example, the Local Group contains about 100 galaxies, with $70 \%$ of them being dwarf galaxies (McConnachie 2012). Dwarf galaxies in the Local Group are dominated by dark matter (Mateo 1998) and a study by Oh et al. (2011) showed for a sample of dwarf galaxies from THINGS (The Hi Nearby Galaxy Survey) that the dark matter mass to total mass ratio in their sample is about 0.7 . If this can be generalised to all of the dwarf galaxies, they may account for a significant mass fraction of the total mass in the Universe.

Naively one could think that considering their small sizes and young chemical ages due to their low metallicities, dwarf galaxies are simple systems that are easy to study. In fact, the star formation and chemical enrichment histories in these galaxies are very complex, and are triggered and sustained by as-yet unknown mechanisms. The wide variety of observed dwarf galaxy populations, starformation activities, chemical enrichment histories and ISM conditions, combined with the fact that observing them is challenging because of their low luminosities, makes it difficult to build a general case for these galaxies. This is illustrated already by the numerous morphological types that have been defined to classify the dwarf galaxy population.

## Morphological classification

Indeed dwarf galaxies can be separated into different families. Of course the classification of dwarf galaxies is still a debated subject. Some classes overlap with some others and the properties of the galaxies within a class are not unique. Moreover the morphology of dwarf galaxies strongly depends on the component we are looking at, e.g., spiral distribution of the atomic gas, presence of a stellar bar or irregular structure of the Hir gas. We follow and describe here the classification from Grebel (2001). We refer to Mateo (1998); Kunth \& Östlin (2000); Grebel (2001) for reviews on the various morphological types for dwarf galaxies and their respective properties.

Dwarf Spirals - dS They are small analogues to the "normal" spiral galaxies ( $\mathrm{S} 0, \mathrm{Sa}, \mathrm{Sb}, \mathrm{Sc}$ and Sd galaxies). They are at the high mass end of the dwarf galaxy distribution, and are characterised by a central surface brightness of $\mu_{V} \geq 23$ mag $\operatorname{arcsec}^{-2}$, Hi masses of $M_{H I} \leq 10^{9} M_{\odot}$ and total masses of $M_{\text {tot }} \leq 10^{10} M_{\odot}$. Late-type dwarf spirals are mostly slow-rotators or exhibit solid-body rotation while early-types dwarfs spirals have rotationally supported exponential disks, and are usually more metal- and gas-rich than late types. Dwarf spirals show a slow continuous level of star formation and are found both in clusters and in the field. An example of a dwarf spiral galaxy is shown in Fig. 2.2 left $^{2}$.

Dwarf Irregulars - dIrr As their name suggest, they have an irregular shape in the optical, with scattered bright Hir regions. These galaxies are characterised by $\mu_{V} \leq 23 \mathrm{mag} \operatorname{arcsec}^{-2}$ (usually well described by an exponential profile), $M_{H I} \leq 10^{9} M_{\odot}$ and $M_{\text {tot }} \leq 10^{10} M_{\odot}$. They are gas-rich and the Hi distribution can show a complicated pattern (with clumps and shells). The most massive dIrrs usually have an Hi distribution that is much more extended than their optical size, and are dominated by solid-body rotation. In less massive dIrrs, rotation is not systematically observed and their Hi distribution can be off-centered or can present a ring-like structure (Young \& Lo 1997). They can present high levels of star-formation activity. They are also found either in clusters or in the field. An example of a dwarf irregular galaxy is shown in Fig. 2.2 center ${ }^{3}$.

Blue Compact Dwarfs galaxies - BCDs These galaxies do not have a specific shape but are usually compact, made of several concentrated star clusters and are gas-rich and actively starforming galaxies (i.e., with levels of star formation that can reach those of "normal" giant galaxies).

[^7]

Fig. 2.2. Examples of (left) a dwarf spiral NGC4214, (center) a dwarf irregular IC10, and (right) a blue compact dwarf IZw18.

Various subtypes of BCDs exist, among which are HiI galaxies (with optical spectra reminiscent of Hir regions), blue amorphous galaxies and Wolf-Rayet (WR) galaxies (which present strong signatures indicative of a significant population of WR stars). Their name is due to their blue colours (in B band) arising from the population of young, hot, massive stars recently formed. Both gas and stars show high central concentrations, thus contributing to a strong and bright central surface brightness $\left(\mu_{V} \leq 19 \mathrm{mag} \operatorname{arcsec}^{-2}\right)$. They are characterised by $M_{H I} \leq 10^{9} M_{\odot}$, often exceeding the mass of stars, and $M_{t o t} \leq 10^{10} M_{\odot}$. Their Hi distribution can also be very extended and sometimes decoupled from the galaxies (van Zee et al. 1998; Cannon et al. 2006). Nonetheless solid-body rotation is observed in the central parts of BCDs. Their star formation is mostly episodic through bursts of star formation. Contrary to dS and dIrrs, they are usually isolated, away from galaxy clusters. An example of a BCD galaxy is shown in Fig. 2.2 right ${ }^{4}$.

Dwarf Spheroidals - dSph have a spheroidal shape and are the faintest and least massive galaxies known. Indeed they are characterised by $M_{V} \geq-14 \mathrm{mag}, \mu_{V} \geq 22 \mathrm{mag} \operatorname{arcsec}^{-2}$ with very little central stellar concentration, $M_{H I} \leq 10^{5} M_{\odot}$ and $M_{t o t} \leq 10^{7} M_{\odot}$. They are often almost devoid of detectable neutral gas, and sometimes their low gaseous content is below the amounts expected from stellar mass loss. They do not show any rotation nor any Hir emission. Their star formation rate is very low. They are usually found close to massive galaxies. An example of a dwarf spheroidal galaxy is shown in Fig. 2.3 left.


Fig. 2.3. Examples of (left) a dwarf spheroidal NGC205, (center) a dwarf elliptical M32 and (right) tidal dwarfs from the NGC7252 system. NGC7252 is in the middle of the image and the two dwarf galaxies are at the edges of the two tidal streams.

[^8]Dwarf Ellipticals - dE have spherical or elliptical shapes, with compact central stellar densities. They are characterised by a high central surface brightness but are gas-poor and not very massive: $M_{V} \geq-17 \mathrm{mag}, \mu_{V} \leq 21 \mathrm{mag} \operatorname{arcsec}^{-2}, M_{H I} \leq 10^{8} M_{\odot}$ and $M_{t o t} \leq 10^{9} M_{\odot}$. As in the case of dSph, they are usually not supported by rotation, show little star formation and are also found close to massive galaxies. An example of a dwarf elliptical galaxy is shown in Fig. 2.3 center ${ }^{5}$.

Tidal dwarfs are a special type of dwarf galaxies as they are the remnants of interactions between two massive galaxies (stripped material or mergers). They form from the debris torn out from massive galaxies during these interactions. Contrary to the other families of dwarfs galaxies they can be metal-rich, for their mass, as they possess the metallicity of their parent galaxies (Duc \& Mirabel 1998). Their characteristics such as mass, luminosity, and gas content, thus mostly depend on the conditions of the interaction and on the properties of their progenitors. An example of an interacting system with two tidal dwarf galaxies is shown in Fig. 2.3 right $^{6}$.

Table 2.1. General characteristics for the various morphological types of dwarf galaxies (except tidal dwarfs). The Milky Way has been added on the last column for comparison.

|  | dS | dIrr | BCDs | dSph | dE | Milky Way |
| :--- | :---: | :---: | :---: | :---: | :---: | :---: |
| $M_{V}[\mathrm{mag}]$ | $\geq-18$ | $\geq-16$ | $\geq-18$ | $\geq-14$ | $\geq-17$ | -21 |
| $\mu_{V}\left[\mathrm{mag} \operatorname{arcsec}{ }^{-2}\right]$ | $\geq 23$ | $\leq 23$ | $\leq 19$ | $\geq 22$ | $\leq 21$ | - |
| $M_{H I}\left[M_{\odot}\right]$ | $\leq 10^{9}$ | $\leq 10^{9}$ | $\leq 10^{9}$ | $\leq 10^{5}$ | $\leq 10^{8}$ | $4.5 \times 10^{9}$ |
| $M_{t o t}\left[M_{\odot}\right]$ | spiral | irregular | compact | elliptical | elliptical | barred spiral |
| Shape | extended | extended | small | small | small | extended |
|  | rich | rich | rich | poor | poor | rich |
| Gas Content | solid-body | solid-body | solid-body | no | no | yes |
| Rotation | slow, cont. | episodic/cont. | episodic and bursty | low, cont. | low, cont. | cont. |
| Star Formation | cluster/field | cluster/field | field | cluster | cluster | - |
| Location |  |  |  |  |  | $10^{10}$ |

Note: cont. $=$ continuous
How these different families of dwarf galaxies are related or not is still an open question. For example Thuan (1985) suggested that BCDs and dIrrs are the same type of galaxy, only differing by their present star-formation rates, and can evolve into dE. On the contrary, James (1994) claimed that dIrrs are a fundamentally different population from BCDs that are in fact dwarf ellipticals undergoing a burst of star formation. Zhao et al. (2013) separated BCDs into irregularand elliptical-types and suggested that irregular- and (some) elliptical-type BCDs might be related and at different stages of galaxy evolution and/or have different progenitors. Additionally, very late-type dwarf spirals may be in a transition phase to dwarf irregulars, and very-low mass dIrrs may be evolving towards dwarf spheroidals (Grebel 2001). Some dIrrs and dSph may also be tidal dwarfs that formed early-on. The question of an evolutionary trend between the different families of dwarf galaxies is thus far from being solved.

In Table 2.1, we summarise the main characteristics for the various dwarf morphological types (except the tidal dwarfs). Our sample of dwarf galaxies is composed of gas-rich dwarf galaxies, mainly BCDs and star-forming dIrrs. Thus, in the next Section, we will see in more detail how the ISM of these gas-rich, star-forming dwarf galaxies differs from more metal-rich environments and we will leave aside the gas-poor types (dEs and dSph galaxies).

[^9]
### 2.2.2 The peculiar ISM of dwarf galaxies

Due to their low metallicity, the ISM of these dwarf galaxies is quite different from the ISM of more massive galaxies, and we review here the main differences in terms of gas and dust properties.

## Gas in low-metallicity galaxies

Star formation - Dwarf galaxies usually have gas densities below the threshold value for star formation (determined by Kennicutt 1989), thus one might conclude that their star-formation activity is very poor. As we saw in Section 2.2 .1 this is certainly not the case for BCDs as well as dIrrs that can be actively forming stars, with star formation being concentrated in one or several knots. Star formation can be either continuous or episodic with bursts of star formation (every $\sim 100 \mathrm{Myr}$ ) between two more quiescent periods of slow continuous star formation (Searle \& Sargent 1972). Without gas inflow, the star-formation activity can cease at anytime, whenever all of the gas has been consumed. Gas-poor dwarfs may thus be galaxies which have already converted most of their gas reservoir into stars. BCDs are the typical environments where episodic star formation can be found. These episodes are usually violent (up to $10 \mathrm{M}_{\odot} \mathrm{yr}^{-1}$, Thuan 1983) and thus short, only a few Myr, as the gas reservoir is consumed extremely rapidly to sustain such a high star-formation rate.

Because of this intense star-formation activity, most dIrrs and BCDs are dominated by a young stellar population concentrated in star-forming complexes, super star clusters (SSCs) or massive O-B associations; embedded in a more extended older stellar population (e.g., Hodge et al. 1991). This older population is most of the time composed of red horizontal branch stars and can be thus younger than globular clusters in the Galaxy. The high star-formation activity in BCDs generates a high number of young and massive stars such as WR stars. When observed in a galaxy, those stars are usually the sign of intense star formation over the past 10 Myr . The presence of such young stars (usually with a higher relative fraction compared to the older stellar population) combined to a lower metallicity of the ISM results in a harder galaxy-wide radiation field (Campbell et al. 1986; Galliano et al. 2003, 2005; Madden et al. 2006). The hardness of the radiation field is defined by the proportion of hard photons, and hence by the shape of the stellar spectrum in the UV. Figure 2.4 shows the UV-visible part of the ISRF of the Galaxy compared to the ISRF of four low-metallicity galaxies, which indeed show a higher level of UV emission.

The hardness of the radiation field can be measured using emission from highly ionised species in the optical (e.g.,OiII $\lambda 4363 / \mathrm{H} \beta$, Nir $\lambda 6584 / \mathrm{H} \alpha$, see Kewley et al. 2001) or in the IR domains (e.g., Neiri $\lambda 15.56 \mu \mathrm{~m} /$ Neir $\lambda 12.81 \mu \mathrm{~m}$, Madden et al. 2006). The radiation field becomes softer as stars evolve off the main sequence after a few Myr (e.g., Levesque et al. 2010). As metallicity increases, the metals present in the atmospheres of massive stars can absorb much of the hard UV photons produced by the star, thus decreasing the hardness of the radiation field in the Hir region. As we will see in the next paragraphs, this harder ISRF at low metallicities will have important consequences on the gas and dust properties.

Atomic gas - The absolute Hi content in low-metallicity galaxies is relatively low compared to moremetal rich environments (see Table 2.1). However, for gas-rich dwarf galaxies, the atomic gas mass fraction is quite high (e.g., $\geq 30 \%$ of the total baryonic mass for BCDs, Thuan \& Martin 1981), compared to that found in more massive galaxies (a percent, or fractions of a percent, see Table 2.1 for our Galaxy). These galaxies thus have a large gas reservoir that has not yet been converted into stars. The Hi distribution is usually extended in some dwarf galaxies and goes beyond the optical radius (in terms of $R_{25}$ ) of these galaxies (see Tully et al. 1978; Thuan \& Martin 1981; Thuan et al. 2004; Begum \& Chengalur 2005a; Pustilnik \& Martin 2007; Hunter et al. 2011). This extended Hi content shows a complex structure that can be a remnant from the formation of the dwarf galaxy


Fig. 2.4. Synthesised ISRF for four low-metallicity galaxies from Galliano et al. (2005), compared to the ISRF of the Galaxy (from Mathis et al. 1983)
or sign of external perturbations (i.e., past or present interactions with a more massive companion). This extended Hı might be seen as a large available reservoir of gas slowly fuelling the centre of the galaxy for star formation, although a direct link with star formation is still debated (Hunter 1997; de Blok \& Walter 2006).

Hi can also trace the dark matter content in galaxies, via the rotation curves. Indeed Hi observations provide information on the rotation velocities of the gas. Using rotation curve models, one can reproduce the gas and stellar radial profiles and infer the mass of the halo using the rotation curve and mass models (usually with the profile from Navarro et al. 1996). Hi observations revealed that dwarf galaxies contain a larger fraction of dark matter, more concentrated in the halo, than more massive galaxies (Brinks \& Taylor 1994).

Molecular gas - Molecular gas, through CO observations, is particularly difficult to detect in dwarf galaxies (i.e., Taylor et al. 1998; Leroy et al. 2009; Schruba et al. 2012), despite their intense starformation activity and large reservoirs of Hi gas. This poses the question of what fuels the star formation in these galaxies. This "CO-deficit" in dwarf galaxies can be interpreted in two ways: either the molecular gas content is indeed very low in dwarf galaxies, but then difficult to reconcile with the necessary amount of molecular gas required to fuel star formation; or CO becomes a poor tracer of $\mathrm{H}_{2}$ at low metallicities, and there may exist a significant amount of $\mathrm{H}_{2}$ not traced by CO. Indeed in low-metallicity ISM, the transparency of the ISM due to lower dust abundances, combined with the increased hardness of the ISRF, allows harder FUV photons to penetrate deeper into the cloud and photodissociate CO, and other molecules that are not efficient at self-shielding (because they are not abundant enough). The transition from Ci to CO occurs thus deeper in the molecular cloud. Additionally, $\mathrm{H}_{2}$ is efficient at self-shielding and can survive in regions where CO has already been photodissociated, resulting in a higher proportion of $\mathrm{H}_{2}$ residing in the CII - CI region, outside the smaller CO core (see Fig. 2.5). This molecular gas not traced by CO has been called the "CO-free" gas (or "CO-dark" gas). Its presence was first suggested in low-metallicity galaxies via observations of [CII] at $158 \mu \mathrm{~m}$ in the Large Magellanic Cloud (Poglitsch et al. 1995; Israel 1997) and IC10 (Madden et al. 1997). Its existence has been confirmed in our Galaxy with two independent measurements: $\gamma$-rays a few years ago (Grenier et al. 2005) and Planck observations in the submm $/ \mathrm{mm}$ domain more recently (Planck Collaboration et al. 2011a). Theoretical work
has shown that the "CO-free" gas effect is strongly controlled by the extinction and should have an increased importance as the metallicity decreases (Wolfire et al. 2010; Krumholz \& Gnedin 2011; Glover \& Clark 2012). In low-metallicity galaxies, recent Herschel and Planck observations indeed confirm that the molecular gas content appears more important that the one traced by CO alone (Roman-Duval et al. 2010; Galliano et al. 2011).


Fig. 2.5. Illustration of the impact of metallicity and extinction on the structure of the molecular clouds, inspired from Wolfire et al. (2010). The harder FUV photons penetrate deeper into the cloud at low metallicity, increasing the relative size of the region where $\mathrm{H}_{2}$ is co-existing with Cir and Ci (yellow zone), compared to the region where both $\mathrm{H}_{2}$ and CO are present (red zone). The blue zone corresponds to the atomic phase, where no $\mathrm{H}_{2}$ is present.

Thus the relation between the total amount of $\mathrm{H}_{2}$ and the observed CO is more ambiguous at low metallicities. Several studies have tried to determined the metallicity dependence of the $X_{C O}$ conversion factor (Wilson 1995; Israel 1997; Taylor et al. 1998; Boselli et al. 2002; Leroy et al. 2011; Schruba et al. 2012). For example, from a sample of 16 dwarf galaxies, Schruba et al. (2012) found a $X_{C O}$ scaling with $(\mathrm{O} / \mathrm{H})^{-2}$.

## Dust in low-metallicity galaxies

Dwarf galaxies have absolute dust abundances lower than those of more metal-rich galaxies, resulting in lower UV-attenuation and a porous ISM. This can be interpreted as a direct consequence of the low metallicity of these galaxies: less metals are available in the ISM to form dust. We saw in the previous section that this has an important influence on the molecular cloud structure. Moreover, despite their low metal content, for their mass, dwarf galaxies possess non-negligible amounts of dust with properties differing from their metal-rich counterparts, in extinction as well as in emission.

Extinction curve - As we saw in Section 1.2.2, the extinction curve depends on the nature and on the properties of the dust present along the line-of-sight. It can thus vary from one galaxy to another, depending on the physical conditions of the ISM. We illustrate the impact of lower metallicity on the extinction curve in Fig. 2.6, by comparing the Milky Way and two low-metallicity dwarf galaxies: the Large and Small Magellanic Clouds (LMC, Z $\sim 1 / 2 Z_{\odot}$ and SMC, $\mathrm{Z} \sim 1 / 5 Z_{\odot}$ ). The three extinction curves are quite similar in the optical range but differ in the UV: the "UV-bump" at 2175 $\AA$ is much less pronounced in the LMC than in the Milky Way and is absent in the SMC; and the FUV slopes also differ. Weingartner \& Draine (2001) reproduced these observed extinction curves in the Magellanic Clouds theoretically with an approximate mixture of silicate and carbonaceous grains. The absence of the $2175 \AA$ feature in the SMC was explained by a deficit in very small
carbonaceous grains (including PAHs) with sizes below $0.02 \mu \mathrm{~m}$.


Fig. 2.6. Interstellar extinction curves for the Milky Way ( $R_{V}=3.1$ ), the LMC and the SMC (adapted from Gordon et al. 2003)

Warmer dust - Dust has been known to be warmer in dwarf galaxies since the first IRAS observations of high $60 / 100 \mu \mathrm{~m}$ flux ratios, compared to that observed for more metal-rich environments (e.g., Helou 1986; Hunter et al. 1989; Sauvage et al. 1990; Melisse \& Israel 1994). These results were then confirmed with Spitzer (Rosenberg et al. 2006; Cannon et al. 2006; Walter et al. 2007; Engelbracht et al. 2008). The SEDs in some low-metallicity star-forming dwarf galaxies indeed peak at shorter wavelengths, sometimes well below $100 \mu \mathrm{~m}$, and have a steeply rising MIR continuum, representative of an overall warmer dust component (Galliano et al. 2003, 2005; Galametz et al. 2009). In comparison, the peak of the SED is around $100-200 \mu \mathrm{~m}$ for more metal-rich galaxies (see Fig. 1.16). This overall warmer dust is a consequence of the harder and more intense ISRF interacting with the ISM of dwarf galaxies (e.g., Madden et al. 2006).

Weak MIR aromatic features - PAHs are often barely detected in these galaxies (e.g., Sauvage et al. 1990; Madden 2000; Wu et al. 2006; Hunter \& Kaufman 2007; Smith et al. 2007), if at all when the metallicity drops to a level of $\sim 1 / 3-1 / 5 Z_{\odot}$ (Engelbracht et al. 2005). The paucity of dust allows the harder UV photons to travel deeper into the ISM and destroy PAH molecules by photoevaporation or photodissociation (Galliano et al. 2003, 2005; Madden et al. 2006; Wu et al. 2006). The dearth of PAH features in dwarf galaxies has also been explained by the destruction of the molecules by SN shocks (O'Halloran et al. 2006) or by a delayed carbon injection in the ISM by asymptotic giant branch (AGB) stars (Galliano et al. 2008).

Grain size distribution - Starbursting dwarf galaxies sometimes show a higher relative abundance of small grains compared to our Galaxy. Mass can indeed be transferred from big grains to smaller grains because of fragmentation and erosion of the big grains by shock waves produced by the numerous SN in these starbursting dwarf galaxies (Jones et al. 1996; Lisenfeld et al. 2002; Galliano et al. 2003, 2005). As these low-metallicity environments are devoid of PAHs grains, too small to survive in such extreme environments, the very small grains are thus the most numerous grains in the ISM. We saw in Section 1.3 that the PAHs were normally the main actors of the gas heating in PDRs around Hir regions. The dearth of PAHs in low-metallicity environments may impact the thermal balance of the gas phase, as the very small grains become the available agents for gas
heating via the photoeletric effect.
Submillimetre excess - An excess emission, unaccountable by usual SED models, is appearing in the FIR to submm/millimetre (mm) domain for some dwarf galaxies (Galliano et al. 2003, 2005; Galametz et al. 2009; Bot et al. 2010; Grossi et al. 2010; Dale et al. 2012). An excess emission has also been observed in our Galaxy with COBE (Reach et al. 1995) but with an intensity less pronounced compared to that found in low-metallicity systems. Dumke et al. (2004); Bendo et al. (2006); Zhu et al. (2009) found a submm excess in some low-metallicity spiral galaxies as well. This submm excess has been one of the main sources of uncertainty in dust modelling in dwarf galaxies for the past few years, especially on the dust mass parameter. Several explanations have already been proposed to investigate the origin of this excess, although not completely satisfactory :

1. An additional very cold dust component: Galliano et al. (2003, 2005), Galametz et al. (2009, 2010 , 2011) modelled the submm excess they detected in their metal-poor galaxies with a very cold dust (VCD) component, residing in very dense clumps. They added to their SED models an extra modified blackbody with a submm emissivity index of 1 and a low dust temperature ( $\sim 10 \mathrm{~K}$ ). Their additional component could explain the break observed in the submm domain in some of their SEDs but it may lead to very low gas-to-dust mass ratios, considering the observed gas mass, compared to those expected from chemical evolution and from the amount of available metals in the ISM. Moreover, Galliano et al. (2011) observed that the submm excess in the LMC is more pronounced in diffuse regions, inconsistent with the assumption of very dense clumps containing this VCD component.
2. An additional spinning dust component: Several studies have shown that fast rotating very small dust grains from ionized gas regions in many galaxies were producing centimetre (cm) radio emission (Ferrara \& Dettmar 1994; Draine \& Hensley 2012). As shown by recent studies (Ysard \& Verstraete 2010; Ysard et al. 2012), the peak of the "spinning" dust emission depends on many parameters such as the radiation field intensity, the dust size distribution, dipole moment distribution, physical parameters of the gas phase, etc. This hypothesis was tested to explain the submm-to-cm excess by Bot et al. (2010), Israel et al. (2010) and Planck Collaboration et al. (2011b) in the Magellanic Clouds. The spinning dust model seems sufficient for the $\mathrm{mm} / \mathrm{cm}$ excess but another effect is required to explain the submm/mm excess. Indeed, the usual spinning dust models do not normally produce much emission in the submm domain but rather at longer wavelengths (as illustrated in Murphy et al. (2010) for NGC 6946). Moreover, PAHs have been assumed to be the carriers of this spinning dust emission (as shown in Draine \& Lazarian 1998) and this may seem contradictory with the weakness of the observed PAH features in low-metallicity galaxies (Engelbracht et al. 2005; Madden et al. 2006).
3. Changes in the optical properties of the grains:

- via a change in the dust grain composition: Welty et al. (2002); Serra Díaz-Cano \& Jones (2008) suggested amorphous carbon as a more realistic ISM dust analogue in lieu of graphite. Amorphous carbon grains are more emissive than graphite grains: the emissivity index, $\beta$, is 1.7 instead of 2.0 (see Sections 1.2 .3 , and Chapter 8). Thus using amorphous carbon instead of graphite in SED models would result in a flatter submm slope and may therefore be sufficient to explain the submm excess (Galametz et al. 2013), and suppress the need to invoke any extra dust component. Additionally, the increased emissivity for the amorphous carbon grains also implies a lower dust mass as less dust is required to account for the same luminosity, and resulting in dust masses and gas-to-dust mass ratios more coherent with those expected from chemical evolution (Galametz et al.

2010; Meixner et al. 2010; O’Halloran et al. 2010; Galliano et al. 2011; Galametz et al. 2013).

- via grain coagulation: Paradis et al. (2009) showed that the FIR-submm excess in dense molecular clouds of the Galaxy could be explained by fractal aggregation of amorphous individual grains which induces changes in the dust submm optical properties. This grain coagulation effect had already been suggested in the past by Bazell \& Dwek (1990) and Stepnik et al. (2001).
- via a temperature-dependent emissivity index: Meny et al. (2007) proposed a new model for FIR/submm dust emission based on the physical properties of disordered matter. They consider the interaction of electromagnetic waves with the acoustic oscillations in a disordered charge distribution (DCD) and a distribution of low energy two level tunnelling states (TLS). This interaction results in an emission spectrum strongly dependent on the temperature and in an enhanced submm/mm emission compared to more classical models. The emissivity index is no longer constant over this wavelength range. This model has been successfully applied by Bot et al. (2010) to explain the mm excess in the LMC. However, when applied to the SMC, DCD/TLS effects alone do not reproduce the excess well (Bot et al. 2010).
- via the addition of magnetic grains: Recent work by Draine \& Hensley (2012) on the SMC focuses on magnetic grains as a possible source of submm excess. They consider nanoparticles of metallic iron, magnetite and maghemite that could be free fliers in the ISM or inclusions on larger dust grains. Magnetic grains indeed have an enhanced absorption cross section at submm wavelengths and part of the submm excess could be due to thermal emission from magnetic grain material. They show that a combination of a normal dust mixture (amorphous silicates and carbonaceous grains), spinning dust and magnetic dust could account for the observed SED in the submm/mm range of the SMC.

We will come back to the submm excess later in Chapter 8 to see how Herschel and new submm data can bring new constraints on this excess, and look at the impact of using amorphous carbon instead of graphite grains.

### 2.2.3 Motivations for dwarf galaxies studies

Besides being the most abundant galaxy population in the Universe, dwarf galaxies pose a number of interesting issues with potentially important consequences on numerous domains in cosmological and extragalactic studies. We illustrate here a few of these issues at different spatial and time scales.

## Large-scale structures and galaxy formation

Dwarf galaxies are interesting objects for cosmological studies for several reasons. As they are dominated by dark matter, they provide constraints on the dark matter distribution predicted by cosmological simulations. Currently these simulations predict an increasing number of small size dark matter halos in which dwarf galaxies reside at the present epoch. However the number of observed dwarf galaxies is way too low compared to the theoretical predictions (Moore et al. 1999). Possible explanations for this discrepancy are that these halos do not contain any visible matter and are thus undetectable (Simon \& Geha 2007), or that these halos host very low surface-brightness dwarf galaxies too difficult to detect. This is known as "the missing satellite problem".

In terms of galaxy formation, dwarf galaxies have long been thought to be building blocks of the larger, more evolved systems of the present-day Universe (Dekel \& Silk 1986). However evidence
for downsizing (Cowie et al. 1996), where star formation occurs first in the massive systems before proceeding to lower mass objects in the local Universe, is challenging this scenario.

## Approaching the primordial Universe

Because of their low metallicity, dwarf galaxies can be considered to be chemically young objects that went through only a few episodes of star formation and that are thus at an early stage of their evolution. In this picture, they can be seen as analogues to primordial environments present in the early Universe. Those systems are very difficult (if not impossible) to observe with the current instrumentation and dwarf galaxies provide interesting local probes of ISM conditions that may resemble those of the early Universe. However the dwarf galaxies we observe today are not pristine environments as the early Universe galaxies are. Indeed, most of the time they also harbour an intermediate-age or old stellar population, signs of a complex star-formation history. This poses the interesting question of the nature and existence of the link between dwarf galaxies and primordial environments beyond chemical youth.

Dwarf galaxies also offer the possibility to approach primordial helium abundances (e.g., Pagel et al. 1992; Izotov et al. 1994, 1997; Olive et al. 1997; Peimbert et al. 2002; Luridiana et al. 2003; Steigman 2005; Izotov et al. 2007, and many others). These abundances in turn provide useful constraints on critical cosmological parameters such as the baryonic mass density of the Universe, the lifetime of the neutron and the number of neutrino families. Dwarf galaxies, and especially BCDs, are ideal laboratories to constrain the helium primordial abundance to the necessary precision. As they are chemically very young, they contain very little helium manufactured by stars and most of the helium present in the ISM of these galaxies is thus primordial helium, produced during the Big Bang nucleosynthesis.

## Galaxy evolution

The external evolution of a galaxy is driven by the environment, as opposed to the internal evolution, driven by chemical enrichment (see next paragraph). This external evolution is controlled by inflows, outflows, interaction with companions or mergers. Simulations show that the more massive galaxies, to reach their current mass and size, must have experienced merging events from smaller systems in the past. As the scenario of dwarf galaxies playing the role of building blocks of more massive galaxies does not seem realistic today, we can wonder if these two main populations of galaxies: dwarfs and "normal" massive galaxies, are linked or if they are two different populations. The continuous evolution of the ISM properties between the dwarfs and massive galaxies seems to point towards the first hypothesis, even if this link is not clear.

## Accessing and understanding the chemical evolution history

As we saw in Section 1.2.7, the ISM matter-life cycle implies that the metallicity of a galaxy increases as the galaxy evolves through several cycles of star formation, building up its stellar mass. Thus we should expect some kind of relation between stellar mass and metallicity, providing us with crucial information on galaxy evolution and on the physical processes governing it. This relation has indeed been observed first locally for a small sample of dwarf galaxies by Lequeux et al. (1979) and extended to a sample of $\sim 50,000$ local star-forming galaxies by Tremonti et al. (2004). The relation between mass and metallicity holds over 3 orders of magnitude, down to $M_{*}=10^{8.5}$ and a factor of 10 in metallicity. The mass-metallicity relation, or $M-Z$ relation, is steep for $M_{*} \leq 10^{10.5}$ and flattens out at larger stellar masses. Tremonti et al. (2004) explained this flattening with chemical evolution models by invoking efficient galactic-scale winds that remove metals from low mass galaxies with shallow potential wells. Using hydrodynamical simulations, Brooks et al. (2007) have shown that in
addition to gas mass loss due to SN feedback, a mass-dependent star-formation efficiency is needed to explain the $M-Z$ relation at low masses: low-mass galaxies with low star-formation efficiencies have not yet converted a lot of their gas reservoirs into stars and are therefore less metal-rich. Alternatively, a lower proportion of high-mass stars in low-mass galaxies, due to low levels of star formation, and thus a slower chemical enrichment has also been proposed (Köppen et al. 2007). The $M-Z$ relation has also been observed at high redshifts (Savaglio et al. 2005; Erb et al. 2006; Maiolino et al. 2008; Mannucci et al. 2009; Cresci et al. 2012) and the relation has been seen evolving until $\mathrm{z} \sim 3$. Thus, investigating the population of low-mass and low-metallicity galaxies, i.e., dwarf galaxies, may provide important clues to the origin of the $M-Z$ relation.

However, metal enrichment is in fact a very complex process and depends on external and internal processes occurring during galaxy evolution. Indeed, the gas phase abundance can be effected by pristine, metal-poor, gas inflows, that will dilute the ISM and decrease the metallicity of the galaxy; or by outflows driven by stellar feedback. Elements are also processed by gas and dust in the ISM as we saw in Chapter 1: elements are depleted from the gas phase when dust is formed and returned to the gas phase when dust is destroyed (see Section 1.2.6). The gas-to-dust mass ratio (G/D) links the amount of metals locked up in dust and in the gas phase and is thus a powerful tracer of the evolutionary stage of the galaxy. Investigation of the relation between G/D and metallicity can thus place important constraints on the physical processes governing galaxy evolution and more specifically on chemical evolution models and their main parameters such as SN rate, initial mass function (IMF), etc. Dwarf galaxies are thus good test cases for these models at low metallicities.

We believe that part of the answer to these issues lies in the understanding of the ISM processes occurring in low-metallicity dwarf galaxies. Throughout Section 2.2 .2 we saw that the ISM of dwarf galaxies present notable peculiarities and differences compared to the ISM of more-metal rich environments, and that several of them are not fully understood yet.

Having reliable probes and diagnostics of the properties of the ISM phases in low-metallicity galaxies is necessary to start tackling these issues. A complete and well-sampled wavelength coverage of the SED is also needed to disentangle the effects of various parameters, such as metallicity, star-formation activity, morphology etc., on the ISM properties and on the processes regulating them. The advent of the Herschel Space Observatory at the very beginning of my PhD, provided new constraints on the FIR-to-submm domain of the SED. Before Herschel, the wavelength ranges accessible to the scientific community with IRAS, ISO, Spitzer, AKARI and ground-based facilities (suffering from contamination by Earth's atmosphere) were indeed poorly sampling the bulk of the dust emission beyond $\sim 160 \mu \mathrm{~m}$ in the FIR-submm and direct observations of the most important gas cooling lines in the FIR were missing. Moreover, the sensitivities reached by these instruments enabled accurate studies only for the brightest and highest metallicity dwarf galaxies. With its unprecedented sensitivity and wavelength coverage, Herschel now provides a unique opportunity to study dust and gas in the lowest-metallicity galaxies, enabling more systematic studies of these environments over a larger range of metallicities.

My work focuses especially on the dust in dwarf galaxies, and on the impact of metallicity on the dust properties. The previous discussion has demonstrated the importance and the need for such a study. The first part of my work (Chapters 3 to 5 ) is dedicated to the study of the cold dust component uncovered by Herschel in low-metallicity environments. In these Chapters, we only focus on new Herschel data (presented in Chapter 4) and interpret it with the aid of a modified blackbody model (Chapter 5). The second part of my work (Chapters 6 to 9 ) broadens our view of dust in these low-metallicity systems by looking at the total dust component across the whole IR-to-submm range (data presented in Chatper 6). The SEDs are modelled with a semi-empirical model, presented in Chapter 7 and the results are then analysed and discussed in Chapter 8. We
especially focus on the variation of the dust properties with metallicity and address the submm excess issue in this chapter. The last chapter of this thesis incorporates the gas into our dust picture and is dedicated to the variations of the G/D with metallicity. Implications for chemical evolution scenarios are discussed with the aid of three chemical evolution models.

I base my work on a survey of low-metallicity galaxies observed with Herschel. I also include another Herschel survey, containing more metal-rich environments to broaden my metallicity range towards higher metallicities and study the variations dust properties from metal-poor to metal-rich systems. The uniqueness of this work is that the dust properties are derived here in a systematic way for a large number of galaxies $(\sim 110)$, with a wavelength range as broad as possible, among which more than half are dwarf galaxies, and $\sim 35 \%$ have $\mathrm{Z} \leq 1 / 5 Z_{\odot}$. In most of the studies mentioned previously, only single systems were studied or small samples containing 10-20 lowmetallicity galaxies, usually with a limited number of galaxies with $\mathrm{Z} \leq 1 / 5 Z_{\odot}$. We saw in Section 2.2.1 that dwarf galaxies are unique systems and that it is very difficult to build a general case for these galaxies. The main challenge of my work is to treat all of the galaxies the same way, to be able to systematically compare their dust properties, without neglecting their individual specificities.

## Part II

## The cold dust in low-metallicity ISMs revealed by Herschel

## Chapter 3

## The Herschel Space Observatory

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In this Chapter, I present the Herschel mission and the two instruments on which most of my work is based. A detailed description of the two Herschel surveys I use for my study is given in the last two sections.

### 3.1 The mission

### 3.1.1 Description

The Herschel Space Observatory was approved by ESA in 1993 as the fourth "cornerstone" mission in the ESA "Horizon 2000" long-term plan. It was successfully launched on the $14^{\text {th }}$ of May 2009, together with Planck, and operated for about 4 years, providing the community with more than 25,000 hours of science data. It operated autonomously far away from Earth on an orbit around the $2^{\text {nd }}$ Lagrangian point of the Sun-Earth/Moon system. Initially named FIRST (Far InfraRed and Submillimetre space Telescope) at the beginning of the mission, it was renamed after the discoverer of infrared light, William Herschel, for the $200^{t h}$ anniversary of this discovery in 2000.

Herschel was designed, following Spitzer, to offer a higher spatial resolution and sensitivity compared to the previous space missions and to explore the relatively unknown FIR/submm wavelength range, where dust emits the bulk of its energy (see Chapter 1). Among the widely-ranging scientific targets of the mission, we can find stars in the early-phase of formation in their birth molecular cloud in the Galaxy, ISMs of local galaxies or high-redshift star-forming galaxies (up to $\mathrm{z} \sim 5$ ). The science objectives of Herschel thus cover all the possible time and spatial scales: from local feedback of evolved stars in the ISM of very nearby galaxies to star-formation history of the universe and galaxy evolution over cosmological times. Herschel was dedicated to the study of the ISM and the cold Universe in its most general sense. In our case, we look at local ISM properties of nearby galaxies and study how dust properties vary in environments of metal content lower than in our Galaxy.


Fig. 3.1. The Herschel Space Observatory, seen on the "warm" side (left) and on the "cold" side (right). The various components of the spacecraft are labelled on the schematic view in the centre.

The mission consists of a spacecraft, equipped with three instruments cooled down to 0.3 K by liquid superfluid helium inside a cryostat. The spacecraft is composed of the payload module carrying the three instruments, the telescope, the service module, and the sunshield/sunshade (Fig. $3.1^{1}$ ).

[^10]

Fig. 3.2. (left) The payload module (from Pilbratt et al. (2010)). (right) Schematic view of the payload module where all of the components have been labelled. The three instruments are highlighted in orange (adapted from the Herschel Observers' Manual).

Autonomous observations according to the timelines, accurate pointings, on-board data handling, and communication with the ground are handled by the spacecraft. The telescope has a typical Cassegrain optical design and a diameter of 3.5 m , making it the largest mirror, that does not need a deployable structure, ever launched in space and reaches unprecedented sensitivities.

Table 3.1. Main capabilities of the three instruments on board Herschel. The first rows, marked with "P" describe the photometers capabilities. The last rows, marked with "S" show the spectrometers capabilities. In bold, the two instruments used in my work: PACS and SPIRE photometers.

|  |  | PACS | SPIRE | HIFI |
| :---: | :---: | :---: | :---: | :---: |
| P | Range | $60-210 \mu \mathrm{~m} 70$ or $110 \mu \mathrm{~m}$ and $160 \mu \mathrm{~m}$ (simultaneous) | $\begin{aligned} & 200-670 \mu \mathrm{~m} \mathrm{250,} 350,500 \mu \mathrm{~m} \\ & \text { (simultaneous) } \end{aligned}$ |  |
|  | FOV | $1.75{ }^{\prime} \times 3.5^{\prime}$ | $4^{\prime} \times 8^{\prime}$ (unfilled) |  |
|  | $\begin{aligned} & \text { Sensitivity } \\ & (5 \sigma / 1 \mathrm{hr}) \\ & \hline \end{aligned}$ | $\begin{aligned} & 5 \mathrm{mJy}(70 / 110 \mu \mathrm{~m}), \\ & 10 \mathrm{mJy}(160 \mu \mathrm{~m}) \\ & \hline \end{aligned}$ | Confusion limit: $5.8 \mathrm{mJy}(250 \mu \mathrm{~m})$, <br> $6.3 \mathrm{mJy}(350 \mu \mathrm{~m}), 6.8 \mathrm{mJy}(500 \mu \mathrm{~m})$ |  |
| S | Range | 55-210 $\mu \mathrm{m}$ | 194-313 and 303-671 $\mu \mathrm{m}$ | $\begin{aligned} & \hline \hline 157-213 \text { and } 240-625 \\ & \mu \mathrm{~m} \text { (with gap) } \end{aligned}$ |
|  | FOV | $47^{\prime \prime} \times 47^{\prime \prime}$ | $2.0^{\prime}$ (unvignetted) | $\begin{aligned} & 39^{\prime \prime}(488 \mathrm{GHz}) \\ & 13^{\prime \prime}(1408 \mathrm{GHz}) \end{aligned}$ |
|  | $\begin{aligned} & \text { Sensitivity } \\ & (5 \sigma / 1 \mathrm{hr}) \end{aligned}$ | $\begin{aligned} & 2 \times 10^{-18} \mathrm{Wm}^{-2}\left(130 \mu \mathrm{~m}, 1^{\text {st }} \text { order }\right), \\ & 5 \times 10^{-18}\left(70 \mu \mathrm{~m}, 3^{\text {rd }} \text { order }\right) \\ & \text { Continuum: } 100 \mathrm{mJy}\left(1^{\text {st }} \text { order }\right), \\ & 250 \mathrm{mJy}\left(3^{\text {rd }} \text { order }\right) \end{aligned}$ | $\begin{aligned} & 1.0-2.2 \times 10^{-17} \mathrm{Wm}^{-2} \text { (high res.) } \\ & 40-88 \mathrm{mJy} \text { (low res.) } \end{aligned}$ | $\begin{aligned} & \text { "A few" mK (Band 1a) } \\ & \text { to } 100 \mathrm{mK} \text { (Band } 7 \mathrm{~b}), \\ & 1 \sigma / 1 \mathrm{hr} \end{aligned}$ |
|  | Resolution | 900-2100 (1st order, 102-210 $\mu \mathrm{m}$ ), 1800-3000 (2nd order, $72-98 \mu \mathrm{~m}$ ), 2600-5400 (3rd order, $55-72 \mu \mathrm{~m}$ ) | 20-1000 | $1000-10^{7}$ |

Note: Table adapted from Tables 3.1 and 3.2 of the Herschel Observers' Manual.
If we zoom into the payload module, we find the three instruments on board Herschel the Photodetector Array Camera and Spectrometer (PACS - Poglitsch et al. 2010), the Spectral and Photometric Imaging REceiver (SPIRE - Griffin et al. 2010) both imaging photometers and medium resolution spectrometers, and the Heterodyne Instrument for the Far Infrared (HIFI - de Graauw
et al. 2010), a very high resolution heterodyne spectrometer (Fig. 3.2). The PACS and SPIRE photometers in combination cover a 60 to $670 \mu \mathrm{~m}$ range, and the three spectrometers together cover 55 to $670 \mu \mathrm{~m}$. The basic informations on the three instruments can be found in Table 3.1.

### 3.1.2 The novelty of Herschel

The era of infrared space astronomy began with the InfraRed Astronomical Satellite (IRAS) in 1983. Since then, several other IR telescopes have been launched including the Infrared Space Observatory (ISO), the Spitzer Space Telescope, the Wide-field Infrared Survey Explorer (WISE) and AKARI. The wavelength ranges and sensitivities covered by Spitzer, ISO and IRAS do not sample the cold dust component of the dust SED beyond $160 \mu \mathrm{~m}$. Some ground-based telescopes such as $J C M T, ~ A P E X, S E S T$, IRAM could detect the cold dust beyond $160 \mu \mathrm{~m}$, but because of sensitivity limitations, accurate measurements of the photometry could only be obtained for the brightest and highest metallicity dwarf galaxies. With unprecedented sensitivity and spatial resolution, Herschel is helping to fill this gap in spectral range and complete our view of dust in the widest range of galaxy types and distances, and, in particular, constraining the cold dust contribution. Table 3.2 summarises various capabilities for the cited space telescopes. Fig. 3.3 illustrates the improvement in spatial resolution brought by Herschel on the nearby dwarf galaxy NGC 1705. The two knots are well separated in both Herschel/PACS images whereas they are blurred in Spitzer/MIPS images and indistinguishable at $160 \mu \mathrm{~m}$.

Table 3.2. Comparison of the infrared telescopes capabilities.

|  | Launch | Wavelength <br> $(\mu \mathrm{m})$ | Spatial resolution <br> at $100 \mu \mathrm{~m}\left({ }^{\prime \prime}\right)$ | Sensitivity <br> at $100 \mu \mathrm{~m}(\mathrm{mJy})$ |
| :--- | :---: | :---: | :---: | :---: |
| IRAS | 1983 | $12-100$ | 150 | $200^{a}$ |
| ISO | 1995 | $2-240$ | 90 | $10^{a}$ |
| Spitzer | 2003 | $3-180$ | $38($ at $160 \mu \mathrm{~m})$ | $1^{b}$ |
| AKARI | 2006 | $1.7-180$ | $39($ at $90 \mu \mathrm{~m})$ | $3.2^{c}$ |
| WISE | 2006 | $2.5-25$ | - | - |
| Herschel | 2009 | $55-670$ | 12 | 1 |
| Planck | 2009 | $350-10^{4}$ | - | - |

${ }^{a}$ : The sensitivities are $1 \sigma$ values for 1 hour integration time in the $100 \mu \mathrm{~m}$ broadband filter.
${ }^{b}$ : The sensitivities are $1 \sigma$ values for 1 hour integration time in the $160 \mu \mathrm{~m}$ broadband filter.
${ }^{c}$ : The sensitivities are $1 \sigma$ values for 2 second integration time in the $90 \mu \mathrm{~m}$ broadband filter.

### 3.2 The PACS and SPIRE photometers

### 3.2.1 PACS

## Description

The PACS photometer performs dual band imaging over a field of view (FOV) of $1.75^{\prime} \times 3.5^{\prime}$, with full sampling of the telescope point spread function (PSF). Two bands can be observed simultaneously: either 70 and $160 \mu \mathrm{~m}$ or 100 and $160 \mu \mathrm{~m}$. The light is spread between the blue ( 70 or $100 \mu \mathrm{~m}$ ) and the red $(160 \mu \mathrm{~m})$ bolometers and is collected by the bolometers arrays in the focal plane (Fig.


Fig. 3.3. Observations of the galaxy NGC1705 with MIPS and PACS: Illustration of the improvement in spatial resolution between Spitzer/MIPS (top row) and Herschel/PACS (bottom row) at $70 \mu \mathrm{~m}$ (left column) and $160 \mu \mathrm{~m}$ (right column) (from Madden et al. 2013). The two knots are well separated at PACS resolutions. The full width at half maximum (FWHM) of the PACS beams for the 70 and $160 \mu \mathrm{~m}$ bands are $5.6^{\prime \prime}$ and $11.4^{\prime \prime}$ while for the MIPS bands the FWHM values are $18^{\prime \prime}$ and $38^{\prime \prime}$. The colour scale is in MJy/sr.
$3.4^{2}$ ). The pointing accuracy has been measured to $2^{\prime \prime}$ at the $1 \sigma$ level on pointed observations. The photometer characteristics are presented in Table 3.3.


Fig. 3.4. (left) Optical layout of the PACS instrument. The photometer optics are on the top left of the picture (framed by the black rectangles). (right) Bolometer matrices of the PACS blue bolometer array. The total blue bolometer array assembly is $64 \times 32$ pixels and is composed of $4 \times 2$ monolithic matrices of $16 \times 16$ pixels, as shown here. In a similar way, the red bolometer array assembly is $16 x 32$ pixels ( 2 x 1 monolithic matrices).

[^11]The PACS PSF (see Fig. 3.5 for the PSF at $100 \mu \mathrm{~m}^{3}$ ), which represents the way the light from a point source is spread onto the detectors, has been measured out to $1000^{\prime \prime}$, using version 7 of the PACS photometric calibration tree available in the Herschel Interactive Processing Environment (HIPE, Ott 2010). This means that a small contribution of the flux from a point source can be found as far as $1000^{\prime \prime}$ away from the centre of the source. The encircled energy fraction (eef) tables, given by HIPE, quantify the fraction of the total flux density contained in a given aperture on the PSFs (inverse of the aperture correction) and are presented in Fig. 3.5. This will have important consequences for determining accurate photometry of our sources (see Chapter 4).


Fig. 3.5. (left) The PACS photometer PSF at $100 \mu \mathrm{~m}$, displayed with a linear scale up to $10 \%$ of the peak. Note the tri-lobe pattern of the PSF, due to an imperfection on the mirror due to the secondary mirror tripod. (centre and right) PACS Encircled Energy Fractions as a function of radius. The right handside is a zoom between 0.8 and 1.0. The dashed line corresponds to the $100 \%$ line and the dotted line to the $95 \%$ line.

The filter transmission curves are shown in Fig. 3.9 and can be extracted from HIPE. PACS has been calibrated on 5 stars for use as primary calibrators and on asteroids plus fainter stars as secondary calibrators. The calibration accuracy is $5 \%$ in each band and is dominated by the calibration model uncertainty.

What is observed by the bolometers, $\bar{F}_{S}$, is directly proportional to the source flux density, $\mathrm{F}_{S}(\nu)$, weighted by the filter transmission, hereafter called the Relative Spectral Response Function, $\mathrm{R}(\nu)$, (RSRF, as in SPIRE Observers' Manual). $\bar{F}_{S}$ is given by:

$$
\begin{equation*}
\bar{F}_{S}=\frac{\int F_{S}(\nu) R(\nu) \mathrm{d} \nu}{\int R(\nu) \mathrm{d} \nu} \tag{3.1}
\end{equation*}
$$

To obtain a monochromatic flux density, the pipeline has to assume a shape for the spectrum of the source and a standard frequency, $\nu_{0}$. For PACS, the standard frequencies are the frequencies corresponding to $\lambda_{0}=\{70,100,160\} \mu \mathrm{m}$. Here the spectral convention is $\nu \times F_{\nu}=$ constant or $\mathrm{F}_{S}(\nu)=\mathrm{F}_{S}\left(\nu_{0}\right) \times\left(\nu_{0} / \nu\right)$, also used for PACS calibration. Importing this in Eq. 3.1 gives:

$$
\begin{equation*}
\bar{F}_{S}=F_{S}\left(\nu_{0}\right) \nu_{0} \frac{\int R(\nu) / \nu \mathrm{d} \nu}{\int R(\nu) \mathrm{d} \nu} \tag{3.2}
\end{equation*}
$$

Inverting Eq. 3.2 will give the monochromatic flux density, i.e., the output of the pipeline, $\mathrm{F}_{S}\left(\nu_{0}\right)$, as a function of what is observed by the bolometers, $\bar{F}_{S}$ :

[^12]\[

$$
\begin{equation*}
F_{S}\left(\nu_{0}\right)=\bar{F}_{S} K_{4} \tag{3.3}
\end{equation*}
$$

\]

where $\mathrm{K}_{4}$ is:

$$
\begin{equation*}
K_{4}=\frac{1}{\nu_{0}} \frac{\int R(\nu) \mathrm{d} \nu}{\int R(\nu) / \nu \mathrm{d} \nu} \tag{3.4}
\end{equation*}
$$

However, there is no particular reason for the source to have a spectral shape $\mathrm{F}_{S}(\nu) \propto \nu^{-1}$. Thus we need to correct the output of the pipeline to the real spectral shape of the source: this is the colour correction.

If the source "true" spectral shape, $\mathrm{F}_{S}^{\prime}(\nu)$, is unknown, we can derive $\mathrm{F}_{S}^{\prime}\left(\nu_{0}\right)$ via:

$$
\begin{equation*}
F_{S}^{\prime}\left(\nu_{0}\right)=F_{S}\left(\nu_{0}\right) \frac{\bar{F}^{\prime} S}{\bar{F}_{S}} \tag{3.5}
\end{equation*}
$$

This transforms into:

$$
\begin{equation*}
F_{S}^{\prime}\left(\nu_{0}\right)=\frac{\int F_{S}^{\prime}(\nu) R(\nu) \mathrm{d} \nu}{\int\left(F_{S}(\nu) / F_{S}\left(\nu_{0}\right)\right) R(\nu) \mathrm{d} \nu} \tag{3.6}
\end{equation*}
$$

Using $\mathrm{F}_{S}(\nu)=\mathrm{F}_{S}\left(\nu_{0}\right) \times\left(\nu_{0} / \nu\right)$, we transform Eq. 3.6 into:

$$
\begin{equation*}
F_{S}^{\prime}\left(\nu_{0}\right)=\frac{\int F_{S}^{\prime}(\nu) R(\nu) \mathrm{d} \nu}{\int\left(\nu_{0} / \nu\right) R(\nu) \mathrm{d} \nu} \tag{3.7}
\end{equation*}
$$

## Observing modes

There are three different observing modes for the PACS photometer: a point source photometry chop-nod mode, a scan map technique for small or large fields, and another scan map technique within the PACS/SPIRE parallel mode. Here we describe shortly the second one as it is the one used for our observations. The map in scan map mode is made by moving the telescope on the sky along parallel lines, called scan legs, at a constant speed (Fig. 3.6). The scan leg length depends on the size of the desired map, and the scan speed for our observations is $20^{\prime \prime} / \mathrm{s}$. To obtain a smoother map, a better PSF reconstruction and to remove the striping effects due to the $1 / f$ noise, it is recommended to do two scan maps with two orthogonal scan directions for one source. For point sources, a mini-scan map mode is preferred over the chop-nod mode, as it increases the point source sensitivity in all bands. It is a special case of the scan map mode with $3^{\prime}$ scan legs and where the scanning is done along the two diagonals of the bolometer detector.


Fig. 3.6. (left) Example of a scan map with 6 scan line legs (from the PACS Observers' Manual). (right) Coverage maps of a scan map and a mini-scan map (from our observations).

### 3.2.2 SPIRE

## Description

The SPIRE photometer observes the three bands simultaneously over a $4^{\prime} \times 8^{\prime}$ field of view. The detector arrays are composed of 139, 88 and 43 bolometers at 250,350 and $500 \mu \mathrm{~m}$ respectively. Several bolometers in each array are centred on the same position on the sky for all three arrays (see blue bolometers in Fig. 3.7 ${ }^{4}$ ).


Fig. 3.7. (left) SPIRE focal plane unit on the photometer side. (right) Schematic view of the bolometers arrays. The blue bolometers are the bolometers centered on the same position on the sky for the 3 arrays. The grey bolometers are the dead bolometers. The $4^{\prime} \times 8^{\prime}$ field of view is delineated by the red dashed rectangle.

The SPIRE photometer characteristics are presented in Table 3.3. The beam areas are given here as they will be needed to convert the flux densities extracted from the map in Jy/beam to Jy (see Chapter 4). Neptune was used as a primary calibrator for the photometer calibration. There are two different SPIRE calibration uncertainties: a systematic uncertainty of $\sim 5 \%$ coming from the models used for Neptune, which is correlated between the three bands, and a random uncertainty of $\sim 2 \%$ coming from the repetitive measurement of the flux densities of Neptune. These two uncertainties are added linearly instead of in quadrature as advised in the SPIRE Observers' Manual, giving an overall $7 \%$ calibration accuracy for the SPIRE photometer ${ }^{5}$.

To treat the signal from the source, the pipeline assumes that the source is a point source, uses the same spectral convention of $\nu \times F_{\nu}=$ constant as PACS, and outputs monochromatic in-beam flux densities at standard frequencies corresponding to 250,350 and $500 \mu \mathrm{~m}$. The reasoning will be the same as for PACS, except that the output will be in units of Jy/beam.

However, for SPIRE, the filter transmission for each band depends on the extent of the source. If the source is extended, the response function has to be weighted by $\lambda^{2}$ to take into account the fact that the beam width varies across the band depending on the wavelength, leading to additionnal power coupled from extended emission. The two response functions for a point source, $\mathrm{R}_{P}(\lambda)$, and an extended source, $\mathrm{R}_{E}(\lambda)$, for each band are shown in Fig. 3.9, with $\mathrm{R}_{E}(\lambda)=\mathrm{R}_{P}(\lambda) \times\left(\lambda_{0} / \lambda\right)^{2}$. The $\mathrm{K}_{4}$ factor, to convert $\bar{F}_{S}$ into $\mathrm{F}_{S}\left(\nu_{0}\right)$ then depends if the source is extended, $\mathrm{K}_{4}=\mathrm{K}_{4 E}$, or not, $\mathrm{K}_{4}=\mathrm{K}_{4}$. As the pipeline does not know if the source is extended or compact, the default will be to assume that the source is a point source, and use $\mathrm{K}_{4 P}$ in Eq. 3.3. Thus to get the

[^13]monochromatic flux density for an extended source, one has to multiply $\mathrm{F}_{S}\left(\nu_{0}\right)$, the output of the pipeline, by $\mathrm{K}_{4 E} / \mathrm{K}_{4 P}$ (see Chaper 4). The $\mathrm{K}_{4 P}$ and $\mathrm{K}_{4 E} / \mathrm{K}_{4 P}$ values are given in Table 3.3.

The colour corrections work the same as for PACS and Eq. 3.7 can be used for SPIRE too. Given the type of source observed by SPIRE, in most cases it could be appropriate to assume that the source spectrum follows a power law with a spectral index, $\alpha_{S}$, i.e., $\mathrm{F}_{S}^{\prime}(\nu)=\mathrm{F}_{S}^{\prime}\left(\nu_{0}\right) \times\left(\nu_{0} / \nu\right)^{\alpha_{S}}$ where $\mathrm{F}_{S}^{\prime}\left(\nu_{0}\right)$ is the source flux density at $\nu_{0}$ for the "true" spectral shape. In this case, Eq. 3.7 transforms into:

$$
\begin{equation*}
F_{S}^{\prime}\left(\nu_{0}\right)=F_{S}\left(\nu_{0}\right) \frac{\nu_{0}}{\nu_{0}^{\alpha_{S}}} \frac{\int R(\nu) / \nu \mathrm{d} \nu}{\int R(\nu) / \nu^{\alpha_{S}} \mathrm{~d} \nu} \tag{3.8}
\end{equation*}
$$

## Observing modes

SPIRE offers three observing modes: a Large area map mode, a Small area map mode and a Point Source Photometry map. We describe briefly here the first two only, as they correspond to the modes used to perform our observations. The Large map mode is used to cover large areas of the sky or extended sources ( $>5^{\prime}$ diameter) and the map is made by scanning the telescope on the sky, at a given scan speed. For our observations, the speed used is the nominal scan speed of $30^{\prime \prime} / \mathrm{s}$. The Small map mode is for sources with diameters $<5^{\prime}$, and the map is made by doing two short cross-scans with the telescope. The difference in coverage is shown in Fig. 3.8 ${ }^{6}$.


Fig. 3.8. Example of a Large (left) and Small (right) scan map coverage with SPIRE.

### 3.3 Low metallicity with Herschel : the Dwarf Galaxy Survey

### 3.3.1 Description

The Dwarf Galaxy Survey (DGS, PI: S. Madden) is one of the three Herschel Guaranteed Time Key Programs of the SPIRE Local Galaxies consortium (SAG2). The two other programs are the Herschel Reference Survey (HRS, PI: S. Eales) and the Very Nearby Galaxy Survey (VNGS, PI: C. Wilson). The DGS aims at studying the gas and dust properties in low-metallicity ISM with Herschel. It is a photometric and spectroscopic survey of 50 dwarf galaxies at FIR and submm wavelengths. For a more detailed description of the general goals of the survey and the source selection process, see the Dwarf Galaxy Survey Overview by Madden et al. (2013). In this work,

[^14]

Fig. 3.9. (top) PACS filter transmissions, $R(\lambda)$. (bottom) SPIRE filter transmissions for point sources, $\mathrm{R}_{P}(\lambda)$, plain lines, and extended sources, $\mathrm{R}_{E}(\lambda)$, dotted lines (see text for details). On both plots, the dashed lines mark the reference wavelength for each band.

Table 3.3. PACS and SPIRE characteristics.

| PACS |  |  |  |  |
| :--- | :---: | :---: | :---: | :---: |
| Band | $\lambda(\mu \mathrm{m})$ | Range $(\mu \mathrm{m})$ | FHWM $\left({ }^{\prime \prime}\right)$ | Calibration accuracy |
| Blue | 70 | $60-85$ | 5.6 | $5 \%$ |
| Green | 100 | $85-125$ | 6.8 | $5 \%$ |
| Red | 160 | $125-210$ | 11.4 | $5 \%$ |


| SPIRE |  |  |  |  |  |  |  |  |
| :--- | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Band | $\lambda(\mu \mathrm{m})$ | Range $(\mu \mathrm{m})$ | FHWM $^{a}\left({ }^{\prime \prime}\right)$ | Beam Areas $^{b}\left({ }^{\prime \prime 2}\right)$ | Calibration accuracy | $\mathrm{K}_{4 P}^{c}$ | $\mathrm{~K}_{4 E}^{c} / \mathrm{K}_{4 P}$ |  |
| PSW | 250 | $200-300$ | 18.2 | 465 | $7 \%$ | 1.0113 | 0.9828 |  |
| PMW | 350 | $270-430$ | 24.9 | 822 | $7 \%$ | 1.0087 | 0.9834 |  |
| PLW | 500 | $375-670$ | 36.3 | 1768 | $7 \%$ | 1.0065 | 0.9710 |  |

${ }^{a}$ : The SPIRE beam FWHM values are given for maps with standard pixel size of $6,10,14^{\prime \prime}$ at $250,350,500$ $\mu \mathrm{m}$.
${ }^{b}$ : SPIRE photometer reference spectrum values: http://herschel.esac.esa.int/twiki/bin/view/Public/ SpirePhotometerBeamProfileAnalysis, September 2012 values. The errors on the beam areas are of $\sim 4 \%$.
${ }^{c}$ : The values have been taken from the SPIRE Observers' Manual, Version 2.4, Section 5.2.8.
we focus on the dust and thus only on the 48 targets for which complete photometry was obtained. Throughout this manuscript, we thus consider that the total sample consists of the 48 photometry targets of the DGS (leaving out the LMC and SMC only done in spectroscopy).

The DGS survey is composed of some well-known Local Group sources (such as IC10, NGC6822, NGC1569) and a selection of sources from several optical line emission surveys including the Hamburg/SAO Survey (i.e., HS0017+1055) and the Second Byurakan Survey (i.e., SBS0335-052). The DGS target sources come from follow-up observations which specifically identified the metal-deficient BCDs from the surveys (e.g., Izotov et al. 1991; Ugryumov et al. 2003; Popescu \& Hopp 2000), as well as the sample of Izotov \& Thuan (1999); Hopkins et al. (2002) which contain some of the better-studied BCDs.

The targets of the DGS span a wide range in various parameters such as morphology, starformation rate, stellar and gas mass and metallicity, with objects among the most metal-poor galaxies in the local Universe (Figs. 3.11 and 3.12). The metallicities range from $12+\log (\mathrm{O} / \mathrm{H})=7.14$ (IZw18, Z $\sim 1 / 40 Z_{\odot}$, Izotov et al. 1999) to 8.43 (He2-10, Z $\sim 0.55 Z_{\odot}$, Kobulnicky et al. 1999). The closest source is NGC 6822 at 500 kpc and the more distant is HS0052+2536 at 191 Mpc (Fig. 3.12). Typical star-formation rates range over 4 orders of magnitude from $0.0007 \mathrm{M}_{\odot} \mathrm{yr}^{-1}$ (UGC4483) to $28 \mathrm{M}_{\odot} \mathrm{yr}^{-1}$ (Haro11), with many of the dwarf galaxies containing super star clusters. The DGS contains mostly objects of Irr type (VIIZw403) along with a few dwarf spiral galaxies (NGC625). Stellar masses range over 4 orders of magnitudes, from $3 \times 10^{6} M_{\odot}(\mathrm{UGC4483})$ to $3 \times 10^{10} M_{\odot}$ (Haro11).

### 3.3.2 Metallicity determination

The most important parameter throughout my study is metallicity. Thus we have to be certain that the metallicities for our samples, are determined in a homogeneous way to avoid any discrepancy or unreliable metallicity trends.

As we saw in Chapter 2, there are several methods for estimating the metallicity of a galaxy. We explore and compare two different methods here: the so-called "direct" method based on the direct measurement of the electron temperature $T_{e}$ (OiII) from the detection of the [OiII] $\lambda 4363$ auroral line, and a "strong-line" method which uses an empirical calibration of metallicity from the [OiI] $\lambda 3727$ and [Oiir] $\lambda \lambda 4959,5007$ strong emission lines. The first method should be more accurate, and closer to the "true" metallicity value, but as the [Oiri] $\lambda 4363$ auroral line is not detected in higher-metallicity environments, we also need to have measurements from a "strong-line" method in order to have a homogeneous set of metallicity values over the largest metallicity range possible. We follow the method and the equations presented in Izotov et al. (2006), hereafter I06, for the "direct" method; and in Pilyugin \& Thuan (2005), hereafter PT05, for the "strong-line" method. The PT05 method uses the $R_{23}$ ratio $^{7}$.

## Direct determination following I06

The $\mathrm{O}^{+}$and $\mathrm{O}^{2+}$ abundances are computed using the following relations:

$$
\begin{array}{r}
12+\log \left(\frac{\mathrm{O}^{+}}{\mathrm{H}^{+}}\right)=\log \frac{\lambda 3727}{H \beta}+5.961+\frac{1.676}{t}-0.40 \log t-0.034 t+\log (1+1.35 x) \\
12+\log \left(\frac{\mathrm{O}^{2+}}{\mathrm{H}^{+}}\right)=\log \frac{\lambda 4959+\lambda 5007}{H \beta}+6.200+\frac{1.251}{t}-0.55 \log t-0.014 t \tag{3.10}
\end{array}
$$

[^15]where $t=10^{-4} \mathrm{~T}_{e}$ and $x=10^{-4} \mathrm{~N}_{e} \mathrm{t}^{-0.5}$. For $\mathrm{O}^{2+}$ in Eq. 3.10, $T_{e}=T_{e}\left(O_{\mathrm{IIII}}\right)$ and for $\mathrm{O}^{+}$in Eq. $3.9, T_{e}=T_{e}\left(O_{\mathrm{II}}\right) . t\left(O_{\mathrm{III}}\right)$ is given by the formula of Aller (1984):
\[

$$
\begin{equation*}
t(O I I I)=\frac{1.432}{\log [(\lambda 4959+\lambda 5007) / \lambda 4363]-\log C_{T}} \tag{3.11}
\end{equation*}
$$

\]

where $\mathrm{C}_{T}$ is given by:

$$
\begin{equation*}
C_{T}=\left(8.44-1.09 t+0.5 t^{2}-0.08 t^{3}\right) \frac{1+0.0004 x}{1.0 .044 x} \tag{3.12}
\end{equation*}
$$

$t$ (OII) can be deduced from $t\left(O_{\text {III }}\right)$ using the relations of Stasińska \& Izotov (2003).
$\mathrm{O}^{3+}$ only represents a small fraction of the total oxygen abundance. Photoionisation models show that $\mathrm{O}^{3+} / \mathrm{O}$ is $\geq 1 \%$ only in the highest excitation regions with $\mathrm{O}^{+} /\left(\mathrm{O}^{+}+\mathrm{O}^{2+}\right) \lesssim 0.1$. In these regions we can approximate $\mathrm{O}^{3+} / \mathrm{H}^{+}$by:

$$
\begin{equation*}
\frac{\mathrm{O}^{3+}}{\mathrm{H}^{+}}=0.5 \frac{\mathrm{He}^{2+}}{\mathrm{He}^{+}+\mathrm{He}^{2+}}\left(\frac{\mathrm{O}^{+}}{\mathrm{H}^{+}}+\frac{\mathrm{O}^{2+}}{\mathrm{H}^{+}}\right) \tag{3.13}
\end{equation*}
$$

using the $\lambda 4471$ line for $\mathrm{He}^{+}$and the $\lambda 4686$ line for $\mathrm{He}^{2+}$. When $\mathrm{O}^{+} /\left(\mathrm{O}^{+}+\mathrm{O}^{2+}\right) \geq 0.1, \mathrm{O}^{3+} / \mathrm{H}^{+} \sim 0$ and we assume we can neglect it. We find that $\mathrm{O}^{3+}$ has to be taken into account in three galaxies only: HS0822+3542, Pox186 and Tol1214-277.

Finally the oxygen abundance is given by summing up all of the ionic abundances:

$$
\begin{equation*}
\frac{\mathrm{O}}{\mathrm{H}}=\frac{\mathrm{O}^{+}}{\mathrm{H}^{+}}+\frac{\mathrm{O}^{2+}}{\mathrm{H}^{+}}+\frac{\mathrm{O}^{3+}}{\mathrm{H}^{+}} \tag{3.14}
\end{equation*}
$$

## Empirical determination following PT05

The relations used to derive the metallicities are the following:

$$
\begin{align*}
12+\log (O / H)_{\text {upper }} & =\frac{R_{23}+726.1+842.2 P+337.5 P^{2}}{85.96+82.76 P+43.98 P^{2}+1.793 R_{23}}  \tag{3.15}\\
12+\log (O / H)_{\text {lower }} & =\frac{R_{23}+106.4+106.8 P-3.40 P^{2}}{17.72+6.60 P+6.95 P^{2}-0.302 R_{23}} \tag{3.16}
\end{align*}
$$

where $P$ is an excitation parameter given by:

$$
\begin{equation*}
P=\frac{[O I I I] \lambda \lambda 4959,5007 / H \beta}{R_{23}}=\frac{[O I I I] \lambda \lambda 4959,5007}{[O I I] \lambda 3727+[O I I I] \lambda \lambda 4959,5007} \tag{3.17}
\end{equation*}
$$

Following Kewley \& Ellison (2008), we use the upper (resp. lower) calibration branch when $\log \left(\left[\mathrm{NIII}^{\mathrm{II}}\right] 6584 /[\mathrm{OII}] \lambda 3727\right)>($ resp. $\leq)-1.2$.

For both methods the errors from the line intensities are propagated to the oxygen abundance using:

$$
\begin{equation*}
\Delta f=\sqrt{\sum_{i=1}^{N}\left(\frac{\partial f}{\partial x_{i}} \Delta x_{i}\right)^{2}} \tag{3.18}
\end{equation*}
$$

for a function f of N variables $\mathrm{x}_{i}, \mathrm{i}=[1 . . \mathrm{N}]$.

## Comparison of the two methods

All of the optical line intensities and their uncertainties are collected from the literature (see references in Table 3.4). For some galaxies, metallicities were computed using the mean $(\mathrm{O} / \mathrm{H})$ of several HII regions. The DGS metallicities for the two methods are listed in Table 3.4. The references listed in Table 3.4 are the references for the original optical line intensities. The comparison between the two methods is shown in Fig. 3.10 and the mean ratio between the two methods is 0.004 dex with an average deviation of 0.10 dex. For $63 \%$ of the galaxies, the two metallicity estimates are within 0.1 dex from each other. And all of the galaxies, except two, have their two metallicity estimates corresponding within 0.2 dex. The two deviant galaxies (Tol0618-402 and UM448) show very high $[\mathrm{OII}] \lambda 3727$ intensities that can boost up their PT05 metallicity compared to their I06 metallicity.


Fig. 3.10. Comparison of the metallicity estimates of the DGS using the methods of I06 and PT05. The unity line is indicated as a dotted line and the mean deviation of 0.1 dex as dashed lines.

For our study, we also use more metal-rich environments (see 3.4), for which only PT05 measurements are available (Moustakas et al. 2010; Kennicutt et al. 2011). Thus from now on, we work with the PT05 metallicity values for the DGS galaxies for consistency. The final metallicity distribution for the DGS sample is presented in Fig. 3.11.

### 3.3.3 Other parameters

The target list is presented in Table 3.5 along with galaxy characteristics:

- (1): Galaxy name.
- (2) and (3): RA (J2000) and DEC (J2000). Positions are from the NASA/IPAC Extragalactic Database (NED).
- (4): Distance, D. Any distance quoted here not referenced to the literature, has been determined using the velocities given in NED and the Mould et al. (2000) model, assuming $\mathrm{H}_{0}=70 \mathrm{~km} \mathrm{~s}^{-1} \mathrm{Mpc}^{-1}$.
- (5): Metallicity, in terms of oxygen abundances: $12+\log (\mathrm{O} / \mathrm{H})$. The values quoted here are the PT05 values, except where noted.
- (6): Absolute magnitude in band B $M_{B}$. They have been determined from the apparent magnitudes, $\mathrm{m}_{B}$, from NED using: $\mathrm{m}_{B}-\mathrm{M}_{B}=5(\log (\mathrm{D})-1)$ where $D$ is the distance in pc.

Table 3.4. Metallicities for the DGS determined following either the I06 or the PT05 method.

| Sources | $12+\log (\mathrm{O} / \mathrm{H}) \mathrm{I} 06$ | $12+\log (\mathrm{O} / \mathrm{H}) \mathrm{PT} 05$ | Ref |
| :---: | :---: | :---: | :---: |
| Haro11 ${ }^{\text {d }}$ | $8.23 \pm 0.03$ | $8.36 \pm 0.01$ | 1 |
| Haro2 ${ }^{\text {a }}$ | - | $8.23 \pm 0.03$ | 2 |
| Haro3 | $8.37 \pm 0.02$ | $8.28 \pm 0.01$ | 3 |
| He2-10 ${ }^{\text {a }}$ | - | $8.43 \pm 0.01$ | 4 |
| HS0017+1055 ${ }^{\text {b }}$ | $7.63 \pm 0.10$ | - | 5 |
| HS0052+2536 ${ }^{\text {b,d }}$ | $8.07 / 8.04 \pm 0.20$ | - | 5 |
| HS0822+3542 | $7.44 \pm 0.06$ | $7.32 \pm 0.03$ | 6 |
| HS1222+3741 | $7.83 \pm 0.03$ | $7.79 \pm 0.01$ | 7 |
| HS1236+3937 | $7.47 \pm 0.04$ | $7.72 \pm 0.01$ | 8 |
| HS1304+3529 | $7.66 \pm 0.04$ | $7.93 \pm 0.01$ | 8 |
| HS1319+3224 | $7.59 \pm 0.04$ | $7.81 \pm 0.01$ | 8 |
| HS1330+3651 | $7.66 \pm 0.04$ | $7.98 \pm 0.01$ | 8 |
| HS1442+4250 | $7.62 \pm 0.02$ | $7.60 \pm 0.01$ |  |
| HS2352+2733 ${ }^{\text {b }}$ | $8.40 \pm 0.20$ | - | 5 |
| $\mathrm{IZw} 18^{\text {d }}$ | $7.11 \pm 0.04$ | $7.14 \pm 0.01$ | 10 |
| IC10 ${ }^{\text {d }}$ | $8.20 \pm 0.06$ | $8.17 \pm 0.03$ | 11 |
| IIZw40 | $8.09 \pm 0.02$ | $8.23 \pm 0.01$ | 12 |
| Mrk1089 ${ }^{\text {c,d }}$ | 8.39/8.21 $\pm 0.12$ | 8.30/8.10 $\pm 0.08$ | 13 |
| Mrk1450 | $8.03 \pm 0.01$ | $7.84 \pm 0.01$ | 14 |
| Mrk153 | $7.83 \pm 0.05$ | $7.86 \pm 0.04$ | 15 |
| Mrk209 | $7.82 \pm 0.01$ | $7.74 \pm 0.01$ | 16 |
| Mrk930 | $8.09 \pm 0.03$ | $8.03 \pm 0.01$ | 17 |
| NGC1140 | $8.27 \pm 0.08$ | $8.38 \pm 0.01$ | 3 |
| NGC1569 | $8.19 \pm 0.03$ | $8.02 \pm 0.02$ | 18 |
| NGC1705 ${ }^{\text {d }}$ | $8.11 \pm 0.13$ | $8.27 \pm 0.11$ | 19 |
| NGC2366 | $7.64 \pm 0.02$ | $7.70 \pm 0.01$ | 20 |
| NGC4214 | $8.20 \pm 0.03$ | $8.26 \pm 0.01$ | 4 |
| NGC4449 ${ }^{\text {a }}$ | - | $8.20 \pm 0.11$ | 21 |
| NGC4861 | $8.05 \pm 0.01$ | $7.89 \pm 0.01$ | 16 |
| NGC5253 | $8.16 \pm 0.03$ | $8.25 \pm 0.02$ | 4 |
| NGC625 ${ }^{\text {d }}$ | $8.18 \pm 0.03$ | $8.22 \pm 0.02$ | 22 |
| NGC6822 ${ }^{\text {d }}$ | $8.11 \pm 0.07$ | $7.96 \pm 0.01$ | 23 |
| Pox186 | $7.75 \pm 0.01$ | $7.70 \pm 0.01$ | 24 |
| SBS0335-052 | $7.29 \pm 0.01$ | $7.25 \pm 0.01$ | 17 |
| SBS1159+545 | $7.46 \pm 0.02$ | $7.44 \pm 0.01$ | 14 |
| SBS1211+540 | $7.69 \pm 0.01$ | $7.58 \pm 0.01$ | 14 |
| SBS1249+493 | $7.72 \pm 0.01$ | $7.68 \pm 0.02$ | 25 |
| SBS1415+437 | $7.61 \pm 0.01$ | $7.55 \pm 0.01$ | 26 |
| SBS1533+574 ${ }^{\text {d }}$ | $8.07 \pm 0.02$ | $8.05 \pm 0.01$ | 16 |
| Tol0618-402 ${ }^{\text {d }}$ | $7.72 \pm 0.01$ | $8.09 \pm 0.01$ | 27 |
| Tol1214-277 | $7.50 \pm 0.03$ | $7.52 \pm 0.01$ | 4 |
| UGC4483 | $7.55 \pm 0.04$ | $7.46 \pm 0.02$ | 28 |
| UGCA $20{ }^{\text {d }}$ | $7.49 \pm 0.01$ | $7.50 \pm 0.01$ | 29 |
| UM133 | $7.70 \pm 0.05$ | $7.82 \pm 0.10$ | 4 |
| UM311 ${ }^{\text {c }}$ | 8.38/8.36 $\pm 0.05$ | $8.38 / 8.36 \pm 0.01$ | 17,30,31 |
| UM448 | $8.03 \pm 0.04$ | $8.32 \pm 0.01$ | 17 |
| UM461 | $7.79 \pm 0.03$ | $7.73 \pm 0.01$ | 15 |
| VIIZw403 | $7.74 \pm 0.01$ | $7.66 \pm 0.01$ | 16 |

Notes:
${ }^{a}$ : For these galaxies, the [OIII] $\lambda 4363$ line was not detected and a direct determination of the metallicity with the I06 method is therefore not possible.
${ }^{b}$ : For these galaxies, no line intensities were available in the litterature. Ugryumov et al. (2003) use the [OIII] 44363 line to get the electron temperature and method from Izotov et al. (1994). I06 is an updated version of the method of Izotov et al. (1994), and thus we chose to quote the results of Ugryumov et al. (2003) in the I06 column.
${ }^{c}$ : These objects are galaxies within compact groups of galaxies or are parts of other galaxies. We chose to quote the value of the metallicity for the galaxy only (on the left) and for the whole group (on the right), which is the mean of all of the metallicities in the group. For Mrk1089, the galaxy is region A-C, and the group is composed of regions A-C, B, E, F1, F2, H from López-Sánchez et al. (2004). For UM311, the galaxy is region 3 following the denomination of Moles et al. (1994), and the group is composed of regions 1-2-3 of Moles et al. (1994) plus NGC450 and UGC807. It was not possible to find a metallicity or lines intensities for UGC807. For HS0052 +2536 the group is composed of HS0052 +2536 and HS0052+2537. For all of the galaxies the group corresponds to the galaxies included in the aperture used for Herschel photometry (see Chapter 4). 66 ${ }^{d}$ : For these galaxies, the final metallicities are derived from the mean $(\mathrm{O} / \mathrm{H})$ of several HII regions.

References for metallicities: (1) Guseva et al. (2012) ; (2) Kong et al. (2002) ; (3) Izotov et al. (2004) ; (4) Kobulnicky et al. (1999) ; (5) Ugryumov et al. (2003) ; (6) Pustilnik et al. (2003) ; (7) Izotov et al. (2007) ; (8) Popescu \& Hopp (2000) ; (9) Guseva et al. (2003a) ; (10) Izotov et al. (1999) ; (11) Magrini \& Gonçalves (2009) ; (12) Guseva et al. (2000) ; (13) López-Sánchez et al. (2004) ; (14) Izotov et al. (1994) ; (15) Izotov et al. (2006) ; (16) Izotov et al. (1997) ; (17) Izotov \& Thuan (1998) ; (18) Kobulnicky \& Skillman (1997) ; (19) Lee \& Skillman (2004a) ; (20) Saviane et al. (2008) ; (21) McCall et al. (1985) ; (22) Skillman et al. (2003) ; (23) Lee et al. (2006) ; (24) Guseva et al. (2007) ; (25) Thuan et al. (1995) ; (26) Guseva et al. (2003b) ; (27) Masegosa et al. (1994) ; (28) van Zee \& Haynes (2006) ; (29) van Zee et al. (1996) ; (30) Moles et al. (1994) ; (31) Pilyugin \& Thuan (2007)

- (7): Galaxy size in terms of $\mathrm{D}_{25}$.
- (8): Stellar mass, $M_{*}$, obtained from the formula of Eskew et al. (2012) using the 3.6 and $4.5 \mu \mathrm{~m}$ IRAC flux densities, $F_{3.6}$ and $F_{4.5}$ derived in Chapter 6:

$$
\begin{equation*}
M_{*}=10^{5.65} F_{3.6}^{2.85} F_{4.5}^{-1.85}\left(\frac{D}{0.05}\right)^{2} \tag{3.19}
\end{equation*}
$$

where $D$ is the distance in Mpc.

- (9): Atomic HI gas mass, $M_{H I}$. If needed, they have been corrected for the distance used here.
- (10): Pre-Herschel total infrared (TIR) luminosity, $L_{T I R}$, determined from the formula of Dale \& Helou (2002), using the 24, 70 and $160 \mu \mathrm{~m}$ Spitzer/MIPS luminosities, $L_{24}, L_{70}$ and $L_{160}$ from Bendo et al. (2012):

$$
\begin{equation*}
L_{T I R}=1.559 \times\left(\nu_{24} L_{24}\right)+0.7686 \times\left(\nu_{70} L_{70}\right)+1.374 \times\left(\nu_{160} L_{160}\right) \tag{3.20}
\end{equation*}
$$

Some galaxies were not observed with one or more of the MIPS bands. When available, we used the IRAS luminosities and the corresponding IRAS formula from Dale \& Helou (2002):

$$
\begin{equation*}
L_{T I R}=2.403 \times\left(\nu_{25} L_{25}\right)+0.7686 \times\left(\nu_{60} L_{60}\right)+1.374 \times\left(\nu_{100} L_{100}\right) \tag{3.21}
\end{equation*}
$$

- (11): Star formation rate, SFR, using the IR-based calibration of Kennicutt (1998):

$$
\begin{equation*}
S F R\left[\mathrm{M}_{\odot} \mathrm{yr}^{-1}\right]=\frac{\mathrm{L}_{\mathrm{TIR}}\left[\mathrm{erg} . \mathrm{s}^{-1}\right]}{2.2 \times 10^{43}}=\frac{\mathrm{L}_{\mathrm{TIR}}\left[\mathrm{~L}_{\odot}\right]}{5.8 \times 10^{9}} \tag{3.22}
\end{equation*}
$$

with $L_{T I R}$ from Column (10). For sources where $L_{T I R}$ was not available, we used $\mathrm{H} \alpha$ or $\mathrm{H} \beta$ luminosities to convert to SFR:

$$
\begin{equation*}
S F R\left[\mathrm{M}_{\odot} \mathrm{yr}^{-1}\right]=\frac{\mathrm{L}_{\mathrm{H} \alpha}\left[\mathrm{erg} . \mathrm{s}^{-1}\right]}{1.26 \times 10^{41}}=\frac{2.85 \times \mathrm{L}_{\mathrm{H} \beta}\left[\mathrm{erg} . \mathrm{s}^{-1}\right]}{1.26 \times 10^{41}} \tag{3.23}
\end{equation*}
$$

The wide range of parameters spanned by the DGS targets is illustrated in Fig. 3.11 and 3.12.

### 3.3.4 Observations with Herschel

The design of the DGS is an attempt to obtain a statistically significant sample of about 9 galaxies in 7 bins over a wide metallicity range: $12+\log (\mathrm{O} / \mathrm{H})$ ranging from $8.4\left(1 / 3 Z_{\odot}\right)$, in bins of


Fig. 3.11. Metallicity distribution of the DGS sample from $12+\log (\mathrm{O} / \mathrm{H})=7.14$ to 8.43 . Solar metallicity is indicated here as a guide to the eye. The pre-Herschel Star Formation Rate ( $S F R$ ) distribution is also indicated by the colour code. They have been converted from $L_{T I R}(S p i t z e r)$ with the Kennicutt (1998) law and are given in Table 3.5. When no IR data was available, $H \alpha$ or $H \beta$ emission lines were used and converted to $S F R$ (Kennicutt 1998) (see Section 3.3.3 for details). The dashed cells indicate that none of these data were available for the galaxy. The most actively star-forming galaxy (in red) corresponds to the starburst luminous infrared galaxy (LIRG) Haro 11.
$12+\log (\mathrm{O} / \mathrm{H})=0.2$ (a factor of 1.4 difference) to provide an accuracy of $\sim 30 \%$. As can be seen from Figure 3.11 , it was not possible to have a uniform sample in terms of metallicity, due to lack of availability or detectability of such extreme low-metallicity sources in the local Universe. We also did not insist on populating our highest metallicity bins, since metal-rich galaxies are plentiful in other surveys, such as the HRS (Boselli et al. 2010b), the Herschel Virgo Cluster Survey (Davies et al. 2012) and the KINGFISH survey (Kennicutt et al. 2011), for example.

The DGS was granted $\sim 230$ hours of observations, part of which were used to observe the sample with the two Herschel imaging photometers: all of them (48) with PACS at 70,100 and $160 \mu \mathrm{~m}$ and 41 with SPIRE at 250,350 and $500 \mu \mathrm{~m}$. Seven sources were not observed with SPIRE because they were predicted to be too faint for SPIRE.

For all of our galaxies, the PACS photometry observations have been done in the PACS scan-map mode, with map sizes varying from $4^{\prime} \times 4^{\prime}$ to $30^{\prime} \times 30^{\prime}$, at a medium scan speed ( $20^{\prime \prime} / \mathrm{s}$ ). Mapping the largest galaxies of the Local Group with the 3 PACS bands required $\sim 10 \mathrm{~h}$ of observing time per galaxy. The SPIRE observations have been made using the SPIRE Large and Small scan-map modes, depending on the source sizes, at the nominal scan speed ( $30^{\prime \prime} / \mathrm{s}$ ). Mapping the 11 larger galaxies takes about 1 h at most per galaxy to reach the desired sensitivities with SPIRE while the small scan mode required about 10 minutes per galaxy. We planned observations to completely cover the star-forming regions and, where possible, much of the HI structure to be able to access the low surface brightness dust in diffuse metal-poor regions. A nominal size of at least $2 \times \mathrm{D}_{25}$ was thus considered for most photometry maps.

The technical details of the DGS observations are summarised in Appendix A.


Fig. 3.12. Properties of the DGS sample: Distributions of distances, $D$, (top left), absolute B magnitudes, $M_{B}$, (top right), stellar masses, $M_{*}$ (bottom left) and HI masses (bottom right). The colour scale represents the range of metallicity values. Galaxies for which the parameter is not available or is an upper limit are not shown here. The number of DGS galaxies on the plot is indicated on the top left of each panel.

### 3.3.5 Ancillary data

The DGS galaxies have been observed with numerous ground-based and space-based observatories (Fig. 3.13). Having the full range of UV to radio observations available opens the door to some very powerful multiphase modelling of the dust and gas, and to a coherent picture of chemical evolution of galaxies. With the vast multi-wavelength Herschel data set, along with the complementary data that exists or is being collected, the physical characteristics of the nebular gas, the photodissociation regions, the neutral gas and stellar properties can be characterized.

The first step in exploiting the Herschel observations has been the construction of the detailed database of existing observations from UV to radio wavelengths. In addition to the new PACS and SPIRE photometry and spectroscopy, the database includes the NIR to mm bands necessary to best characterize the dust properties using existing IRAS data, WISE NIR to MIR measurements, Spitzer MIR to FIR observations, the critical ground-based submm observations: $850 \mu \mathrm{~m}, 870 \mu \mathrm{~m}$ and 1.2 mm data from SCUBA/JCMT or SCUBA-2/JCMT, LABOCA/APEX, MAMBO/IRAM as well as Planck data, which is complementing the longer submm to radio wavelengths. We also include Spitzer IRS spectroscopy to access the atomic, ionized and molecular gas phases. To characterize the stellar activity, 2MASS and GALEX data have also been collected.


Fig. 3.13. Histogram of the number of observed and detected Herschel sources in the DGS program along with the accompanying ancillary data that has been collected. Note that PACS spectroscopy includes 7 individual LMC and SMC sources, thus accounting for the larger number ( 55 sources) in the histogram while the accompanying LMC and SMC Herschel photometry data is in the HERITAGE program.

The ancillary data compiled for the DGS sample to date include the following:

- GALEX maps in FUV and NUV (1528 and 2271£). Observations are collected from the STScI MAST archive (http://galex.stsci.edu/GR6/). These data give us a view into the starformation histories of the galaxies by capturing the direct UV emission from young stars to the most recent star-formation activity (e.g., Hunter et al. 2010). The dust absorption properties in the UV provide additional constraints on the SED modelling of the dust emission properties in the MIR to FIR. GALEX observations exist for all but 9 galaxies of the DGS sample.
- 2MASS observations have been compiled from IRSA (http://irsa.ipac.caltech.edu/) and exist for 34 galaxies in the DGS. The J, H and $\mathrm{K}_{S}$ bands from 2MASS provide direct access to the emission from the oldest stellar population.
- Spitzer has observed all 50 targets of the DGS with various combinations of IRAC bands at $3.6,4.5,5.8$ and $8.0 \mu \mathrm{~m}$ and MIPS bands at 24,70 and $160 \mu \mathrm{~m}$. The IRAC and MIPS data consist of a combination of archival data and the cycle 5 program: Dust Evolution in Low-Metallicity Environments (P.I. F. Galliano; ID: 50550). The MIPS data are reduced and compiled in Bendo et al. (2012) and have been recently publicly released. Spitzer data will enable us to constrain the MIR part of the SEDs and get access, for example, to the PAH features (through the 8 and $24 \mu \mathrm{~m}$ bands) or to the old stellar population (through the first IRAC bands). Furthermore MIPS shares two bands with PACS at 70 and $160 \mu \mathrm{~m}$. This will be very useful to compare the two instruments and confirm the reliability of our PACS results (see Chapter 6). Spitzer IRS spectroscopy, covering 5 to $35 \mu \mathrm{~m}$, exists for all but 5 of the DGS galaxies and have been recently reprocessed (see Chapter 6). The IRS spectroscopy covers the PAHs bands as well as many ionized lines that are relatively bright in dwarf galaxies.
- The sensitivity of $\operatorname{IRAS}$ at $12,25,60$ and $100 \mu \mathrm{~m}$ provides data for about half of the DGS sample (data from the NASA/IPAC IRSA database). The $12 \mu \mathrm{~m}$ observations provide an additional valuable MIR constraint useful for modelling the global SEDs of the DGS galaxies.
- Most of the DGS galaxies have also been detected by WISE at 3.4, 4.6, 12 and $22 \mu \mathrm{~m}$ (data from the NASA/IPAC IRSA database). The $22 \mu \mathrm{~m}$ observations, especially, will give the necessary constraints on the MIR wavelength range of the SED when neither IRAS $25 \mu \mathrm{~m}$ nor Spitzer $24 \mu \mathrm{~m}$ are available.
- Submm observations from ground-based telescopes are crucial to determine the presence and characterization of the submm excess (see Chapter 8). Usually this excess may begin to appear around $450-500 \mu \mathrm{~m}$ compelling observations beyond $500 \mu \mathrm{~m}$. About $30 \%$ of the DGS sample have already been observed in the submm range with either the MAMBO 1.2 mm (IRAM 30 m telescope), LABOCA (APEX), SCUBA(JCMT) or SCUBA-2 (JCMT). After submitting several proposals, we obtained new observations for DGS sources with LABOCA on APEX, presented in Section 6.5.3, to augment the submm observations to characterize and study the submm excess.
- Planck has observed the sky with bands from 30 to 857 GHz . Due to the relatively large Planck beams ( $4.5^{\prime}$ to $33^{\prime}$ ) and limited sensitivity, only a handful of galaxies (eleven) are included in the Planck Early Release Catalog and in the Planck Catalog of Compact Sources.

A summary of the available ancillary data can be found in Table 3.6. Figure 3.13 gives a view of the range of existing Herschel and ancillary data for the DGS sample, and how many galaxies were detected for the different data sets.

### 3.4 Comparing with more metal-rich galaxies: the KINGFISH sample

To enrich our study of the DGS and determine the impact of metallicity on the dust properties, we perform a comparison with galaxies from another Herschel sample: Key Insights on Nearby Galaxies: a Far-Infrared Survey with Herschel (KINGFISH, Kennicutt et al. 2011). KINGFISH is a survey including more metal-rich galaxies and enables us to span a wider metallicity range, notably by populating the high-metallicity end of the metallicity distribution (Fig. 3.14).

The KINGFISH survey contains 61 galaxies: 41 spiral galaxies, 11 early-type galaxies (E and S0) and 9 irregulars (Kennicutt et al. 2011), all observed in the 6 Herschel bands. It has been built with 57 galaxies from the Spitzer Infrared Nearby Galaxies Survey (SINGS, Kennicutt et al. 2003a), to which were added IC0342, NGC2146, NGC 3077 and NGC5457. The metallicities adopted here for most of the KINGFISH sample have been determined the same way as for the DGS in Moustakas et al. (2010), using the method of Pilyugin \& Thuan (2005). The KINGFISH metallicities are from Column 9 from Table 1 of Kennicutt et al. (2011). We have to note also that for 12 galaxies, no Hir region measurement is available in the literature and the metallicities were estimated from the luminosity-metallicity relation. No errors for metallicities are given in Kennicutt et al. (2011) so we adopt a conservative 0.1 dex error for the KINGFISH metallicities. For the four additional sources, IC0342, NGC2146, NGC 3077 and NGC5457, Kennicutt et al. (2011) gives a reference for the metallicity but do not precise if the metallicities were just extracted from the referenced paper or recomputed from original line intensities. Thus for these four sources we recompute the metallicities using the original line intensities from the literature:

- For IC0342, the line ratios were taken from McCall et al. (1985) and we derive a PT05 metallicity of 8.46, close to the value reported by Kennicutt et al. (2011): 8.49.
- For NGC2146, Kennicutt et al. (2011) uses the metallicity determined by Engelbracht et al. (2008) which computed the metallicity using the PT05 method. We thus keep the metallicity given by Kennicutt et al. (2011) for this galaxy.
- For NGC5457, we find a metallicity of $12+\log (\mathrm{O} / \mathrm{H})=8.42$ using the optical line ratios of the 25 regions studied by Kennicutt et al. (2003a). This value is a bit lower than the value reported by Kennicutt et al. (2011): 8.68.
- For NGC3077, no metallicity is reported for the PT05 calibration in Kennicutt et al. (2011). We used McQuade et al. (1995) optical fluxes and derived a PT05 metallicity of $12+\log (\mathrm{O} / \mathrm{H})$ $=8.52$ for this galaxy. The metallicity reported by Kennicutt et al. (2011) using the theoretical calibration of Kobulnicky \& Kewley (2004) is 8.9. The difference between the two metallicity estimates, 0.4 dex , is of the same order than that observed for the other KINGFISH galaxies.

The distances, metallicities and TIR luminosities can be found in Kennicutt et al. (2011). The stellar masses are derived from the IRAC fluxes densities (from Dale et al. 2007) using Eq. 3.19, and most of the HI masses can be found in Draine et al. (2007) (see Chapter 9). The SFR are derived from the TIR luminosities and Eq. 3.22. Fig. 3.14 shows the properties of the KINGFISH sample. This sample spans a 1.2 dex range in metallicity, ranges over 4 orders of magnitudes in stellar mass and 3 orders of magnitude in Hı mass. The DGS distributions for each parameter are overlaid on Fig. 3.14 to facilitate the comparison between the two samples. We see that the two samples complement each other very well in terms of metallicities, stellar and gas masses (Fig. 3.14 and Table 3.4). Indeed the KINGFISH sample brings in more metal-rich galaxies (the median metallicity is 8.4 compared to 7.9 for the DGS) with higher stellar masses (the median $\log \left(M_{*}\right)$ is 10.4 compared to 8.6 for the DGS) and higher gas masses (the median $\log \left(M_{H I}\right)$ is 9.1 compared to 8.6 for the DGS).

Since most of the KINGFISH sources are from SINGS, a large range of ancillary data is available for these galaxies, especially UV to submm photometry (Dale et al. 2005, 2007), including 2MASS, Spitzer, IRAS and SCUBA. For a more detailed description of the available data for the KINGFISH survey, see Kennicutt et al. (2011). The Herschel data is presented in Dale et al. (2012).


Fig. 3.14. Properties of the KINGFISH sample: Distributions of distances, $D$, (top left), metallicities, in terms of $12+\log (\mathrm{O} / \mathrm{H})$, (top right), stellar masses, $M_{*}$ (bottom left) and HI masses (bottom right). The colour scale represents the range of star-formation rate for the top right panel and metallicity for the others. Colour scales are identical as for Fig. 3.11 and 3.12. Galaxies for which the parameter is not available or is an upper limit are not shown here. The number of KINGFISH galaxies on the plot is indicated on the top left of each panel. The DGS distributions for each parameter are indicated in grey.

Table 3.4. Main characteristics of the DGS and KINGFISH parameters.

|  |  | $D$ | $12+\log (\mathrm{O} / \mathrm{H})$ | $M_{B}$ | $\log \left(M_{*}\right)$ | $\log \left(M_{H I}\right)$ | $\log (S F R)$ |
| :--- | :--- | :---: | :---: | :---: | :---: | :---: | :---: |
|  | $[\mathrm{Mpc}]$ |  |  | $\left[\log \left(M_{\odot}\right)\right]$ | $\left[\log \left(M_{\odot}\right)\right]$ | $\left[\log \left(\mathrm{M}_{\odot} \mathrm{yr}^{-1}\right)\right]$ |  |

Table 3.5. Characteristics of the DGS sample.

| Source | $\begin{gathered} \hline \text { RA } \\ (\mathrm{J} 2000) \end{gathered}$ | DEC (J2000) | Distance <br> Mpc (ref) | $12+\underset{(\text { ref })}{\log (\mathrm{O} / \mathrm{H})}$ | $\begin{gathered} \hline M_{B} \\ \text { (ref) } \end{gathered}$ | $\begin{gathered} \hline \text { Size }\left(D_{25}\right) \\ \operatorname{arcmin} \times \operatorname{arcmin} \end{gathered}$ | $\begin{gathered} M_{\star} \\ 10^{8} M_{\odot} \\ \hline \end{gathered}$ | $\begin{gathered} M_{H I} \\ 10^{8} M_{\odot}(\mathrm{ref}) \\ \hline \end{gathered}$ | $\begin{gathered} \hline \hline \log \left(L_{T I R}\right) \\ \log \left(L_{\odot}\right) \\ \hline \end{gathered}$ | $\begin{gathered} S F R \\ \mathrm{M}_{\odot} \mathrm{yr}^{-1} \\ \hline \end{gathered}$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Haro 11 | 00h36m52.7s | -33d33m17.0s | 92.1 (1) | $8.36 \pm 0.01$ (1) | -20.55 (1) | 0.5 'x0.4' | 172.0 | 5.0 (1) | 11.22 | 28.61 |
| Haro 2 | 10 h 32 m 32.0 s | $+54 \mathrm{~d} 24 \mathrm{~m} 02.0 \mathrm{~s}$ | 21.7 (2) | $8.23 \pm 0.03$ (2) | -18.28 (1) | 1.2 'x0.8' | 36.7 | 3.8 (2) | 9.72 | 0.91 |
| Haro 3 | 10 h 45 m 22.4 s | $+55 \mathrm{~d} 57 \mathrm{~m} 37.0 \mathrm{~s}$ | 19.3 (3) | $8.28 \pm 0.01$ (3) | -18.29 (1) | 1.2 'x0.8' | 30.9 | 11.3 (3) | 9.67 | 0.81 |
| He 2-10 | 08 h 36 m 15.1 s | -26d24m34.0s | 8.7 (21) | $8.43 \pm 0.01$ (4) | -17.25 (1) | 1.9 'x1.4' | 31.8 | 3.1 (4) | 9.66 | 0.79 |
| HS 0017+1055 | 00 h 20 m 21.4 s | +11d12m21.0s | 79.1 (3) | $7.63 \pm 0.10^{a}$ (5) | -16.23 (2) | - | 0.69 | $\leq 1.7$ (5) | - | - |
| HS $0052+2536$ | 00 h 54 m 56.4 s | +25d53m08.0s | 191.0 (3) | $8.07 / 8.04 \pm 0.20^{a, b}$ (5) | -19.56 (2) | - | 204.0 | $\leq 482.0$ (6) | - | - |
| HS $0822+3542$ | 08 h 25 m 55.5 s | $+35 \mathrm{~d} 32 \mathrm{~m} 32.0 \mathrm{~s}$ | 11.0 (4) | $7.32 \pm 0.03$ (6) | $-12.40^{d}$ (3) | 0.27 'x0.12' | 0.03 | 0.77 (7) | $\leq 7.04$ | $\leq 0.002$ |
| HS 1222+3741 | 12 h 24 m 36.7 s | $+37 \mathrm{~d} 24 \mathrm{~m} 37.0 \mathrm{~s}$ | 181.7 (3) | $7.79 \pm 0.01$ (7) | -17.90 (4) | $0.12{ }^{\text {'x0 } 0.11 ' ~}$ | 23.7 | - | - | 1.51 |
| HS 1236+3937 | 12 h 39 m 20.2 s | +39d21m05.0s | 86.3 (3) | $7.72 \pm 0.10^{c}$ (8) | -16.00 (4) | $0.34{ }^{\prime} \times 0.09{ }^{\prime}$ | $\leq 14.9$ | - | - | 0.09 |
| HS 1304+3529 | 13 h 06 m 24.2 s | $+35 \mathrm{~d} 13 \mathrm{~m} 43.0 \mathrm{~s}$ | 78.7 (3) | $7.93 \pm 0.10^{c}$ (8) | -17.73 (4) | 0.28 'x0.21' | 3.8 | - | - | 0.50 |
| HS 1319+3224 | 13 h 21 m 19.7 s | +32d08m25.0s | 86.3 (3) | $7.81 \pm 0.10^{c}$ (8) | -15.68 (4) | $0.14{ }^{\prime} \times 0.11$ ' | 1.1 | - | - | - |
| HS 1330+3651 | 13 h 33 m 08.3 s | +36d36m33.0s | 79.7 (3) | $7.98 \pm 0.10^{c}$ (8) | -17.81 (5) | 0.29 'x0.19' | 8.9 | - | - | 0.38 |
| HS 1442+4250 | 14 h 44 m 12.8 s | +42d37m44.0s | 14.4 (3) | $7.60 \pm 0.01$ (9) | -15.15 (4) | 1.13 'x0.26' | - | 3.1 (5) | $\leq 7.59$ | $\leq 0.007$ |
| HS 2352+2733 | 23 h 54 m 56.7 s | +27d49m59.0s | 116.7 (3) | $8.40 \pm 0.20^{a}$ (5) | -17.05 (2) | - | $\leq 1.3$ | $\leq 230.0$ (6) | - | - |
| I Zw 18 | 09h34m02.0s | $+55 \mathrm{~d} 14 \mathrm{~m} 28.0 \mathrm{~s}$ | 18.2 (5) | $7.14 \pm 0.01$ (10) | -15.22 (1) | 0.3 'x0.3' | 0.2 | 1.0 (8) | $\leq 7.74$ | $\leq 0.01$ |
| II Zw 40 | 05h55m42.6s | $+03 \mathrm{~d} 23 \mathrm{~m} 32.0 \mathrm{~s}$ | 12.1 (20) | $8.23 \pm 0.01$ (11) | -14.93 (1) | 0.56 'x0.22' | 4.0 | 5.7 (9) | 9.40 | 0.43 |
| IC 10 | 00 h 20 m 17.3 s | $+59 \mathrm{~d} 18 \mathrm{~m} 14.0 \mathrm{~s}$ | 0.7 (6) | $8.17 \pm 0.03$ (12) | -12.43 (1) | 6.8 'x5.9' | $2.6{ }^{e}$ | 1.1 (10) | $7.80{ }^{\text {f }}$ | 0.01 |
| Mrk 1089 | 05h01m37.7s | -04d15m28.0s | 56.6 (3) | $8.30 / 8.10 \pm 0.08^{b}(13)$ | -20.46 (1) | 0.61 'x0.23' | 192.0 | 210.0 (11) | $10.60^{f}$ | 6.86 |
| Mrk 1450 | 11 h 38 m 35.7 s | +57d52m27.0s | 19.8 (3) | $7.84 \pm 0.01$ (14) | $-13.86^{d}$ (3) | 0.40 'x0.31' | 1.2 | 0.43 (12) | 8.44 | 0.05 |
| Mrk 153 | 10 h 49 m 05.0 s | $+52 \mathrm{~d} 20 \mathrm{~m} 08.0 \mathrm{~s}$ | 40.3 (3) | $7.86 \pm 0.04$ (15) | -18.03 (1) | 0.54 'x0.30' | 11.4 | $\leq 6.9$ (13) | 8.97 | 0.16 |
| Mrk 209 | 12 h 26 m 15.9 s | +48d29m37.0s | 5.8 (7) | $7.74 \pm 0.01$ (16) | -13.67 (1) | $0.8{ }^{\prime} \times 0.6{ }^{\prime}$ | 0.29 | 0.74 (13) | 7.49 | 0.005 |
| Mrk 930 | 23 h 31 m 58.3 s | +28h56m50.0s | 77.8 (3) | $8.03 \pm 0.01$ (17) | - | 0.43 'x0.43' | 27.5 | 31.9 (14) | 10.26 | 3.12 |
| NGC 1140 | 02h54m33.6s | -10d01m40.0s | 20.0 (8) | $8.38 \pm 0.01$ (3) | -18.55 (1) | 1.7 'x0.9' | 44.2 | 35.0 (24) | 9.52 | 0.57 |
| NGC 1569 | 04h30m49.0s | +64d50m53.0s | 3.1 (9) | $8.02 \pm 0.02$ (18) | -15.75 (1) | 3.6 'x1.8' | 8.4 | 1.8 (15) | 9.04 | 0.19 |
| NGC 1705 | 04h54m13.5s | -53d21m40.0s | 5.1 (10) | $8.27 \pm 0.11$ (19) | -15.73 (1) | 1.9 'x1.4' | 2.2 | 0.9 (16) | 7.82 | 0.01 |
| NGC 2366 | 07h28m54.6s | +69d12m57.0s | 3.2 (11) | $7.70 \pm 0.01$ (20) | -16.18 (1) | 8.1 'x3.3' | 1.9 | 7.4 (17) | 8.13 | 0.02 |
| NGC 4214 | 12 h 15 m 39.2 s | $+36 \mathrm{~d} 19 \mathrm{~m} 37.0 \mathrm{~s}$ | 2.9 (12) | $8.26 \pm 0.01$ (4) | -17.11 (1) | 8.5 'x6.6' | 8.4 | 4.1 (15) | 8.73 | 0.09 |
| NGC 4449 | 12 h 28 m 11.1 s | +44d05m37.0s | 4.2 (13) | $8.20 \pm 0.11$ (21) | -17.94 (1) | 6.2 'x4.4' | 26.9 | 11.0 (15) | 9.33 | 0.36 |
| NGC 4861 | 12 h 59 m 02.3 s | +34d51m34.0s | 7.5 (14) | $7.89 \pm 0.01$ (16) | -16.84 (1) | 4.0 'x1.5' | 3.1 | 4.8 (18) | 8.49 | 0.05 |
| NGC 5253 | 13 h 39 m 55.9 s | -31d38m24.0s | 4.0 (12) | $8.25 \pm 0.02$ (4) | -16.87 (1) | 5.0 'x1.9' | 5.9 | 1.6 (19) | 9.14 | 0.24 |
| NGC 625 | 01h35m04.6s | -41d26m10.0s | 3.9 (15) | $8.22 \pm 0.02$ (22) | -16.20 (1) | 5.8 'x1.9' | 5.4 | 1.1 (29) | 8.41 | 0.04 |
| NGC 6822 | 19 h 44 m 57.7 s | -14d48m12.0s | 0.5 (16) | $7.96 \pm 0.01$ (23) | -13.19 (1) | 15.5 'x13.5' | 1.6 | 1.3 (20) | 7.6 | 0.007 |
| Pox 186 | 13 h 25 m 48.6 s | -11d36m38.0s | 18.3 (3) | $7.70 \pm 0.01$ (24) | $-12.17^{d}$ (3) | - | 0.12 | $\leq 0.02$ (21) | - | 0.04 |
| SBS 0335-052 | 03h37m44.0s | -05d02m40.0s | 56.0 (3) | $7.25 \pm 0.01$ (17) | -16.02 (6) | 0.23 'x0.20' | 1.1 | 4.6 (22) | $\leq 9.23$ | $\leq 0.29$ |
| SBS 1159+545 | 12 h 02 m 02.4 s | +54d15m50.0s | 57.0 (3) | $7.44 \pm 0.01$ (14) | -15.08 (7) | 0.20 'x0.10' | 0.70 | $\leq 0.63$ (5) | - | 0.17 |
| SBS 1211+540 | 12 h 14 m 02.5 s | $+53 \mathrm{~d} 45 \mathrm{~m} 17.0 \mathrm{~s}$ | 19.3 (3) | $7.58 \pm 0.01$ (14) | -13.43 (8) | 0.26 'x0.17' | 0.40 | 0.56 (12) | - | 0.02 |
| SBS 1249+493 | 12 h 51 m 52.5 s | +49d03m28.0s | 110.8 (3) | $7.68 \pm 0.02$ (25) | - | 0.10 'x0.10' | 2.7 | 10.0 (23) | - | - |
| SBS 1415+437 | 14 h 17 m 01.4 s | +43d30m05.0s | 13.6 (17) | $7.55 \pm 0.01$ (26) | $-14.99^{d}$ (3) | 0.75 'x0.15' | $0.94{ }^{\text {e }}$ | 4.4 (12) | $\leq 7.78$ | $\leq 0.01$ |
| SBS 1533+574 | 15 h 34 m 13.8 s | $+57 \mathrm{~d} 17 \mathrm{~m} 06.0 \mathrm{~s}$ | 54.2 (3) | $8.05 \pm 0.01$ (16) | $-16.69^{d}$ (3) | 0.67 'x0.67' | 19.9 | 30.0 (14) | $\leq 10.30^{f}$ | $\leq 3.44$ |
| Tol 0618-402 | 06h20m02.5s | -40d18m09.0s | 150.8 (3) | $8.09 \pm 0.01$ (27) | - | $0.4{ }^{\prime} \times 0.3$ ' | 271.0 | - | $\leq 9.91$ | $\leq 1.40$ |
| Tol 1214-277 | 12 h 17 m 17.1 s | -28d02m33.0s | 120.5 (3) | $7.52 \pm 0.01$ (4) | -17.00 (9) | 0.08 'x0.05' | 1.5 | $\leq 3.2$ (5) | $\leq 9.33^{g}$ | $\leq 0.36$ |
| UGC 4483 | 08h37m03.0s | +69d46m31.0s | 3.2 (11) | $7.46 \pm 0.02$ (28) | -12.53 (1) | 1.4 'x0.7' | 0.07 | 0.33 (28) | 6.61 | 0.0007 |
| UGCA 20 | 01h43m14.7s | +19d58m32.0s | 11.0 (18) | $7.50 \pm 0.02$ (29) | -14.43 (2) | 3.1 'x0.8' | $\leq 0.35$ | 0.20 (25) | - | 0.01 |

Table 3.5. Characteristics of the DGS sample (continued).

| Source | RA | DEC | Distance <br> Mpc (ref) | $\overline{\prime \prime} 12+\underset{(\text { ref })}{\log (\mathrm{O} / \mathrm{H})}$ | $\begin{gathered} \hline \hline M_{B} \\ \text { (ref) } \end{gathered}$ | Size $\left(D_{25}\right)$ arcmin $\times$ arcmin | $\begin{gathered} \hline M_{\star} \\ 10^{8} M_{\odot} \end{gathered}$ | $\begin{gathered} M_{H I} \\ 10^{8} M_{\odot}(\mathrm{ref}) \end{gathered}$ | $\begin{gathered} \hline \hline \log \left(L_{T I R}\right) \\ \log \left(L_{\odot}\right) \end{gathered}$ | $\begin{gathered} S F R \\ \mathrm{M}_{\odot} \mathrm{yr}^{-1} \end{gathered}$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| UM 133 | 01h44m41.3s | +04d53m26.0s | 22.7 (3) | $7.82 \pm 0.01$ (4) | $-15.72^{d}$ (3) | 1.0 'x1.0' | 0.65 | 4.0 (26) | - | 0.02 |
| UM 311 | 01h15m34.4s | -00d51m46.0s | 23.5 (3) | $8.38 / 8.36 \pm 0.01^{b}(17,30,31)$ | -19.22 (10) | 0.11 'x0.10' | $103.0^{e}$ | 620.0 (27) | 9.73 | 0.94 |
| UM 448 | 11 h 42 m 12.4 s | +00d20m03.0s | 87.8 (3) | $8.32 \pm 0.01$ (17) | -20.02 (1) | 0.4 'x0.4' | 416.0 | 60.0 (24) | 10.86 | 12.36 |
| UM 461 | 11 h 51 m 33.3 s | -02d22m22.0s | 13.2 (3) | $7.73 \pm 0.01$ (15) | -13.73 (6) | 0.30 'x0.22' | 0.46 | 1.3 (28) | 7.80 | 0.01 |
| VII Zw 403 | 11 h 27 m 59.9 s | +78d59m39.0s | 4.5 (19) | $7.66 \pm 0.01$ (16) | -13.77 (1) | 1.4 'x0.8' | 0.16 | 0.52 (2) | 7.23 | 0.003 |

References for distances : (1) Bergvall et al. (2006) ; (2) Kennicutt et al. (2003a) ; (3) this work, calculated from the redshifts available in NED, the Hubble flow model from Mould et al. (2000) and assuming $\mathrm{H}_{0}=70 \mathrm{~km} \mathrm{~s}^{-1} \mathrm{Mpc}^{-1}$; (4) Pustilnik et al. (2003) ; (5) Aloisi et al. (2007) ; (6) Kim et al. (2009) ; (7) Schulte-Ladbeck et al. (2001) ; (8) Moll et al. (2007); (9) Grocholski et al. (2012) ; (10) Tosi et al. (2001) ; (11) Karachentsev et al. (2002) ; (12) Karachentsev et al. (2004); ; (13) Karachentsev et al. (2003); (14) de Vaucouleurs et al. (1991) ; (15) Cannon et al. (2003) ; (16) Gieren et al. (2006) ; (17) Aloisi et al. (2005) ; (18) Sharina et al. (1996) ; (19) Lynds et al. (1998) ; (20) Bordalo et al. (2009) ; (21)
Tully References for metallicities: (1) Guseva et al. (2012) ; (2) Kong et al. (2002) ; (3) Izotov et al. (2004) ; (4) Kobulnicky et al. (1999) ; (5) Ugryumov et al. (2003) ; (6) Pustilnik et al. (2003) ; (7) Izotov et al. (2007) ; (8) Popescu \& Hopp (2000) ; (9) Guseva et al. (2003a) ; (10) Izotov et al. (1999); ;(11) Guseva et al. (2000) ; (12) Magrini \& Gonçalves (2009); ; (19) Lee \& Skillman (2004b) ; (20) Saviane et al. (2008) ; (21) McCall et al. (1985) ; (22) Skillman et al. (2003) ; (23) Lee et al. (2006) ; (24) Guseva et al. (2007) ; (25) Thuan et al. (1995) ; (26) Guseva et al. (2003b) ; (27) Masegosa et al. (1994) ; (28) van Zee \& Haynes (2006) ; (29) van Zee et al. (1996) ; (30) Moles et al. (1994) ; (31) Pilyugin \& Thuan (2007) References for apparent magnitudes: (1) de Vaucouleurs et al. (1991) ; (2) Ugryumov et al. (2003) ; (3) Gil de Paz \& Madore (2005); (4) Vennik et al. (2000) ; (5) Ugryumov
et al. (2001) ; (6) Maddox et al. (1990) ; (7) SDSS DR4; (8) Paturel et al. (2003) ; (9) Fricke et al. (2001) ; (10) Smoker et al. (2000) et al. (2001) ; (6) Maddox et al. (1990) ; (7) SDSS DR4 ; (8) Paturel et al. (2003) ; (9) Fricke et al. (2001) ; (10) Smoker et al. (2000)
References for HI masses: (1) Cormier et al. (2014) ; (2) Thuan et al. (2004) ; (3) Gordon \& Gottesman (1981) ; (4) Sauvage et al References for HI masses: (1) Cormier et al. (2014) ; (2) Thuan et al. (2004) ; (3) Gordon \& Gottesman (1981) ; (4) Sauvage et al. (1997) ; (5) Pustilnik \& Martin (2007) ; (6)
MacHattie et al. (in prep.) ; (7) Chengalur et al. (2006) ; (8) Lelli et al. (2012) ; (9) Bettoni et al. (2003) ; (10) Huchtmeier \& Richter (1988) ; (11) Williams et al. (1991) ; (12) Huchtmeier et al. (2005) ; (13) Thuan \& Martin (1981) ; (14) Thuan et al. (1999a) ; (15) Walter et al. (2008) ; (16) Meurer et al. (1998) ; (17) Hunter et al. (2011) ; (18) van Eymeren et al. (2009) ; (19) López-Sánchez et al. (2012) ; (20) de Blok \& Walter (2006) ; (21) Begum \& Chengalur (2005b) ; (22) Ekta et al. (2009) ; (23) Pustilnik et al. (2002) ; (24) Davoust \& Contini (2004) ; (25) Paturel et al. (2003) ; (26) Ekta \& Chengalur (2010) ; (27) Moles et al. (1994) ; (28) van Zee et al. (1998) ; (29) Cannon et al. (2004)
$a$ For these galaxies, no line intensities were available in the literature to determine the metallicities using the PT05 method. We use Ugryumov et al. (2003) where the [OIII] 14363
line is used to obtain the electron temperature and the method from Izotov et al. (1994). I06 is an updated version of the method of Izotov et al. (1994).

These objects are galaxies within compact groups of galaxies or are parts of other galaxies. We list the value of the metallicity for the galaxy only ( $1^{\text {st }}$ value quoted) and for the
whole group ( $2^{\text {nd }}$ value quoted), which is the mean of all of the metallicities in the group. For Mrk 1089 , the galaxy is region A-C, and the group is composed of regions A-C, B, E, F1, F2, H from López-Sánchez \& Esteban (2010). For UM311, the galaxy is region 3 following the denomination of Moles et al. (1994), and the group is composed of regions 1-2-3 of Moles et al. (1994) plus NGC450 and UGC807. It was not possible to find a metallicity or lines intensities for UGC807. For HS0052+2536 the group is composed of HS0052+2536
${ }^{c}$ For these galaxies, no line intensity uncertainties, from which the metallicities are determined, are quoted. We assume a conservative value of 0.1 which is the mean of the variation ${ }_{d}$ of the difference between this metallicity calibration, PT05, and the I06 method (see Section 3.3.2). $d$ corrected for Galactic extinction.
${ }^{e}$ computed with synthesised IRAC
$e$ computed with synthesised IRAC fluxes from the SED modelled in Chapter 7.
$f$ IRAS fluxes from NED.
${ }^{g}$ MIPS fluxes from Engelbracht et al. (2008)
Table 3.6. Summary of available Herschel photometry and ancillary data for the DGS galaxies.

| Source | Herschel FIR/submm | $\begin{gathered} \hline \text { GALEX }^{a} \\ \text { NUV/FUV } \end{gathered}$ | $\begin{gathered} \hline \hline \text { 2MASS }{ }^{b} \\ \text { NIR } \end{gathered}$ | $\begin{gathered} \hline \hline \text { IRAS }^{b} \\ \text { IR } \end{gathered}$ | Spitzer $^{C}$ MIR/FIR | $\begin{gathered} \hline \text { WISE }^{b} \\ \text { NIR/MIR } \end{gathered}$ | Ground-Based ${ }^{d}$ submm | $\begin{gathered} \text { Plancke bands }(\mathrm{GHz}) \\ \text { submm } \end{gathered}$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Haro 11 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{f}$ | WISE | LABOCA | - |
| Haro 2 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{g}$ | WISE | SCUBA | - |
| Haro 3 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{f}$ | WISE | - | 857 |
| He 2-10 | PACS - SPIRE |  | JHK | IRAS | IRAC - MIPS - IRS ${ }^{f}$ | WISE | SCUBA MAMBO | 857-545 |
| HS 0017+1055 | PACS - SPIRE | - | JHK | - | IRAC - MIPS24-IRS ${ }^{f}$ | WISE | - | - |
| HS 0052+2536 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS24-IRS ${ }^{f}$ | WISE | LABOCA | - |
| HS 0822+3542 | PACS | GALEX | JHK | - | IRAC - MIPS - IRS ${ }^{f}$ | WISE | - | - |
| HS 1222+3741 | PACS | GALEX | JHK | - | IRAC - MIPS70 - $\mathrm{IRS}^{f}$ | WISE | - | - |
| HS 1236+3937 | PACS - SPIRE | GALEX | - | - | IRAC - $\mathrm{IRS}^{\text {f }}$ | - | - |  |
| HS 1304+3529 | PACS - SPIRE | - | JHK | - | IRAC - MIPS24-IRS ${ }^{f}$ | WISE | - | - |
| HS 1319+3224 | PACS | - | - | - | IRAC - $\mathrm{IRS}^{f}$ | WISE | - | - |
| HS $1330+3651$ | PACS | - | - | - | IRAC - IRS ${ }^{f}$ | WISE | - | - |
| HS 1442+4250 | PACS - SPIRE | GALEX | - | - | IRAC4.5/8 - MIPS - IRS ${ }^{f}$ | WISE | - | - |
| HS 2352+2733 | PACS - SPIRE | GALEX | - | - | MIPS24- IRS ${ }^{f}$ | WISE | - | - |
| I Zw 18 | PACS - SPIRE | GALEX | JHK | - | IRAC - MIPS - IRS $^{f, g}$ | WISE | - | - |
| II Zw 40 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS ${ }^{h}$ - IRS ${ }^{f}$ | WISE | SCUBA - MAMBO | - |
| IC 10 | PACS - SPIRE | - | JHK | IRAS | IRAC - MIPS ${ }^{h}$ - IRS $^{f, g}$ | WISE | SCUBA | 857-545 |
| Mrk 1089 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS24/70- $\mathrm{IRS}^{\text {f,g }}$ | WISE | LABOCA | - |
| Mrk 1450 | PACS - SPIRE | - | JHK | IRAS | IRAC - MIPS - IRS ${ }^{f}$ | WISE | - | - |
| Mrk 153 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS ${ }^{h}$ - IRS ${ }^{f}$ | WISE | - | - |
| Mrk 209 | PACS - SPIRE | GALEX | JHK | - | IRAC4.5/8 - MIPS - IRS $^{f}$ | WISE | - | - |
| Mrk 930 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{f}$ | WISE | LABOCA | - |
| NGC 1140 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{f}$ | WISE | SCUBA - MAMBO | 857 |
| NGC 1569 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{f, g}$ | WISE | SCUBA - MAMBO | - |
| NGC 1705 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{g}$ | WISE | LABOCA | - |
| NGC 2366 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{\text {f,g }}$ | WISE | - | 857 |
| NGC 4214 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{f, g}$ | WISE | SCUBA | 857-545-353-217 |
| NGC 4449 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{g}$ | WISE | SCUBA - MAMBO | 857-545-353-217 |
| NGC 4861 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{\text {f,g }}$ | WISE | - | - |
| NGC 5253 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS ${ }^{h}$ - IRS $^{g}$ | WISE | SCUBA - LABOCA | 857-545-353 |
| NGC 625 | PACS - SPIRE | GALEX | JHK | IRAS | MIPS - IRS ${ }^{f}$ | WISE | LABOCA | 857-545 |
| NGC 6822 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS $^{f, g}$ | WISE | LABOCA | 857 |
| Pox 186 | PACS - SPIRE | - | - | - | IRAC - MIPS24-IRS ${ }^{f}$ | WISE | - | - |
| SBS 0335-052 | PACS - SPIRE | GALEX | JHK | - | IRAC - MIPS - IRS ${ }^{f}$ | WISE | - | - |
| SBS 1159+545 | PACS - SPIRE | - | - | - | IRAC - MIPS24-IRS ${ }^{f}$ | WISE | - | - |
| SBS 1211+540 | PACS - SPIRE | GALEX | - | - | IRAC - MIPS24-IRS ${ }^{f}$ | WISE | - | - |
| SBS 1249+493 | PACS - SPIRE | GALEX | - | - | IRAC - MIPS24-IRS ${ }^{f}$ | WISE | - | - |
| SBS 1415+437 | PACS | GALEX | - | - | IRAC4.5/8-MIPS - IRS ${ }^{f}$ | WISE | - | - |
| SBS 1533+574 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS70 - IRS ${ }^{f}$ | WISE | - | - |
| Tol 0618-402 | PACS | GALEX | JHK | - | IRAC - MIPS - IRS ${ }^{f}$ | WISE | - | - |
| Tol 1214-277 | PACS - SPIRE | GALEX | - | - | IRAC - MIPS ${ }^{h}-$ IRS $^{f}$ | WISE | - | - |
| UGC 4483 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{f}$ | WISE | - | - |
| UGCA 20 | PACS | GALEX | - | - | IRAC - MIPS24-IRS ${ }^{f}$ | WISE | - | - |

Table 3.6. Summary of available Herschel photometry and ancillary data for the DGS galaxies.

| Source | $\begin{gathered} \hline \hline \text { Herschel } \\ \text { FIR/submm } \end{gathered}$ | $\begin{gathered} \hline \hline \text { GALEX }^{a} \\ \text { NUV/FUV } \end{gathered}$ | $\begin{gathered} \hline \hline \text { 2MASS } \\ \text { NIR } \end{gathered}$ | $\begin{gathered} \hline \hline \text { IRAS }^{b} \\ \text { IR } \\ \hline \end{gathered}$ | $\begin{gathered} \hline \hline \text { Spitzer }^{C} \\ \text { MIR/FIR } \end{gathered}$ | $\begin{gathered} \hline \hline \text { WISE }^{b} \\ \text { NIR/MIR } \end{gathered}$ | $\begin{gathered} \text { ground-based } \\ \text { submm } \end{gathered}$ | $\begin{aligned} & \hline \hline \text { Plancke }{ }^{e} \text { bands }(\mathrm{GHz}) \\ & \text { submm } \end{aligned}$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| UM 133 | PACS - SPIRE | GALEX |  |  | IRAC - MIPS24- IRS $^{g}$ | WISE |  |  |
| UM 311 | PACS - SPIRE | GALEX | - | - | IRAC4.5/8-MIPS - IRS $^{f}$ | WISE | LABOCA | 857-545 |
| UM 448 | PACS70/160-SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{f}$ | WISE | SCUBA - LABOCA | - |
| UM 461 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS ${ }^{f}$ | WISE | - | - |
| VII Zw 403 | PACS - SPIRE | GALEX | JHK | IRAS | IRAC - MIPS - IRS $^{f, g}$ | WISE | - | - |

$a$ : Data from STScI MAST archive (http://galex.stsci.edu/GR6/).
b. Data collected from the NASA/IPAC Infrared Science Archive (IRSA).
d. Data collected from the literature: SCUBA/JCMT: 450 and/or $850 \mu \mathrm{~m}$; LABOCA/APEX: $870 \mu \mathrm{~m}$; MAMBO/IRAM: 1.2 mm (e.g., Bolatto et al. 2000; James et al. 2002; Lisenfeld et al. 2002; Böttner et al. 2003; Galliano et al. 2003; Kiuchi et al. 2004; Galliano et al. 2005; Galametz et al. 2009, 2011) or recent follow-up observations (Rémy-Ruyer et al., in preparation).
:- Data collected from the Planck database and the Early Release Source Catalog. Planck bands at 857, 545, 353, 217 and 143 GHz respectively correspond to $350,550,850,1380$
$f:$ Spitzer IRS 1D spectrum available for at least one of the modules: ShortLow, ShortHigh, LongLow, LongHigh. Data is extracted using the database tool, CASSIS (Lebouteiller et al. 2011).
g. Spitzer IR
$h$. While MIPS data are technically available for all MIPS bands, the data from one or two of the bands may have limited usefulness (limited coverage in the case of IC 10; confusion issues in the $160 \mu \mathrm{~m}$ images for Mrk 153 and Tol 1214-277; saturated emission at $24 \mu \mathrm{~m}$ in NGC 5253 ; problems with background subtraction in the $160 \mu \mathrm{~m}$ image of II Zw40 (see Bendo et al. 2012)).

## Chapter 4

## Building the cold dust SEDs

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In this Chapter, we describe the data reduction process followed to produce the final Herschel maps for the DGS sample (Section 4.1). In Sections 4.2 and 4.3, we describe how we obtain the different PACS and SPIRE flux densities, together with their uncertainties (Tables 4.1 and 4.2).

### 4.1 Reducing DGS Herschel data

Herschel data can be reduced using the HIPE environment. During my PhD, HIPE was constantly evolving and several versions of HIPE have been used for the data reduction. We mention each time which version is used for the PACS and SPIRE data reduction ${ }^{1}$.

### 4.1.1 PACS

For the PACS data reduction we use HIPE, with version 7 of the photometric calibration ${ }^{2}$, and a modified version of the pipeline provided by the Herschel Science Center. We describe the main steps of the data reduction that we follow.

The pipeline begins with the Level 0 Products, at a purely instrumental level. All of the auxiliary data (such as "housekeeping" parameters, pointings, etc) is stored as Products. Level 0 also contains the Calibration Tree, needed for flux conversion. Then we perform the usual steps such as flagging the "bad" saturated pixels, converting the signal into $\mathrm{Jy} \mathrm{pix}^{-1}$ and applying flatfield correction. We systematically mask the column 0 of all the matrices (the PACS array is composed of groups of $16 \times 16$ bolometers) to avoid electrical crosstalk issues. We perform second level deglitching to remove all of the glitches, which represent on average $\sim 0.3 \%$ of the data.

After performing all of the above steps we reach the Level 1 stage of data reduction. Note that we still have the bolometer drifts (the so-called " $1 / f^{\prime}$ noise) at this stage of the data reduction. This low-frequency noise originates from two sources: thermal noise, strongly correlated between the bolometers, and uncorrelated non-thermal noise. The method employed to remove the drifts will greatly affect the final reconstructed map (also called Level 2 data). We thus analyse three different map-making methods in order to systematically compare the maps and extract flux densities, to determine if there is an optimized method for each galaxy. The first two map-making methods are provided in HIPE: the PhotProject and the MADmap method. The last method is the Scanamorphos method (Roussel 2013).

The first technique we use for the final reconstruction of the map is РнотРroject. We first remove the $1 / f$ noise (corresponding to data with low spatial frequencies or large scale structures in the map) using a high-pass filter. We then use the РнотProject routine to reproject the data on the sky. The high-pass filtering step is optimum for compact sources but can lead to suppression of extended features (corresponding to low spatial frequencies) in extended sources.

MADmap (Microwave Anisotropy Dataset mapper) produces maximum likelihood maps from the time ordered data (Cantalupo et al. 2010). The main assumption here is that the noise is uncorrelated from pixel to pixel. However, one component of the $1 / f$ noise is strongly correlated from pixel to pixel, as it is due to the thermal drift of the bolometers, and thus not treated by MADmap. Nevertheless, MADmap is more efficient than РнотРroject in reconstructing the extended structures within a map.

[^16]

Fig. 4.1. Scanamorphos (left) and PhotProject (right) images of IC 10 at $70 \mu \mathrm{~m}$ to illustrate how PhotProject tends to clip out the extended features. The colours and spatial scales are the same on both images. Here the diffuse extended emission is best visible on the Scanamorphos map. The comparison of the total flux densities coming from the 2 methods confirms that PhotProject misses the extended emission: in this case, $F_{70}($ Scanamorphos $) / F_{70}($ РhotProject $)=1.5$.

Scanamorphos is another technique specially developed to process scan observations (Roussel 2013). The particularity of Scanamorphos, compared to MADmap, is that no particular noise model is assumed to deal with the low-frequency noise (the $1 / f$ noise). Indeed Scanamorphos takes advantage of the redundancy in the observations, i.e., of the fact that a portion of the sky is observed more than once and by more than one bolometer. The two noise sources contributing to the low-frequency noise are inferred from the redundancy of the data and removed (Roussel 2013). The maps are made using the default parameters. We add the minimap option when reducing data with a field size of the order of 8.4 arcmin . For consistency in the following flux computation, we produce maps with the same pixel sizes for all of the methods: $2^{\prime \prime}, 2^{\prime \prime}$ and $4^{\prime \prime}$ for 70,100 and $160 \mu \mathrm{~m}$ respectively.

## Adapting the map-making method to the source

To determine the best map-making method for each galaxy (summarised in Table 4.1), we compute the flux densities (see Section 4.2.1 for PACS flux extraction) for the three bands for the three methods for each galaxy and compare the photometry for the three different methods. For consistency, we use the same apertures for the three different types of maps.

As mentioned above, the РнотProject method is optimized for compact sources. Indeed, the filtering step partly removes large scale structures in the map. It is not adapted for extended sources as this filtering step can sometimes also remove the large scale structures of our large sources with diffuse extended emission (Figure 4.1), also noted by Aniano et al. (2012) for two extended KINGFISH galaxies. Moreover the source is automatically masked before the high-pass filtering step, and this mask may be too small for extended sources with peculiar morphology, leading to suppression of extended features during the filtering step. Therefore, we decide to take as final, the maps produced by РнотProject for compact sources only.

Some galaxies are not detected in one or several bands. When deriving upper limits on the flux densities for these galaxies, the three methods give very different results. As the "non-detection" criterion is directly linked to the background determination through its contribution to the total flux density and the corresponding uncertainty, we need to choose the method that gives the most reliable background structure. MADmap and Scanamorphos do not have any constraints on
the background values, whereas РнотProject is constrained to an average statistically-null background. Because Scanamorphos does not make any assumptions on the background, sometimes positive residual noise structures can remain in the maps. MADMAP presents features, such as a curved background for some maps, due to an oversimplified treatment of missing data. Again, the РнотProject maps here are used because they are the most constrained as far as the background is concerned. Moreover, when the galaxy is not detected at $160 \mu \mathrm{~m}$ it is usually a compact point source at the other PACS wavelengths. Thus this choice is consistent with the previous choice for compact sources.

For more extended sources, we only consider MADmap and Scanamorphos. As mentioned before, MADMAP maps sometimes present a curved background: the source in the map centre is surrounded by lower background levels than those used in the background aperture for the photometry. This therefore results in a high background leading to an underestimation of the source flux density. Moreover this is not consistent with the assumption of a flat background made for the photometry (see Section 4.2.1). After some tests, to avoid this problem, we decide to use the Scanamorphos maps for the extended sources.

### 4.1.2 SPIRE

Following the same method as in Ciesla et al. (2012) for the Herschel Reference Survey, or in Auld et al. (2013) for the HErschel VIrgo Cluster Survey, the SPIRE maps are processed through HIPE ${ }^{3}$ using a modified version of the available SPIRE pipeline. The steps from the Level 0 to Level 1 are basically the same as in the official version provided by the SPIRE Instrument Control Centre (ICC). The pipeline starts with a first deglitching step, then a time response correction is applied to match the detector timelines to the astronomical pointing timelines. A second deglitching step is then performed as it improves the removal of residual glitches. After an additional time response correction, the flux calibration step is performed, where non-linearity corrections are taken into account. An additional correction is applied to the bolometer timelines to account for the fact that there is a delay in the response of the bolometers to the incoming signal. The temperature drifts of the bolometers are then removed. For this step, the pipeline temperature drift removal is not run. Instead a custom temperature drift correction (BriGAdE, Smith et al. in prep) is applied to the whole observation timeline (rather than to an individual scan-leg). Finally, the Naive mapper is used to construct the final map with pixel sizes of $6,8,12^{\prime \prime}$ for the $250,350,500$ $\mu \mathrm{m}$ band respectively. For galaxies with heavy cirrus contamination, an additional destriping step is performed. A complete description of the data processing step will be given in Smith et al. in prep.

### 4.2 DGS PACS photometry

Throughout this study we focus on the global SED modelling of the samples. Due to the intrinsically faint luminosities of dwarf galaxies, the total integrated photometry for the DGS must be done with great care, paying close attention to the evaluation of uncertainties on the flux densities.

### 4.2.1 Extracting the fluxes

For PACS measurements, we perform aperture photometry, placing an aperture on the source and a background region away from the source to estimate the sky level. As mentioned in Section 3.2.1, the PACS PSFs have been measured out to $1000^{\prime \prime}$ with version 7 of the PACS photometric calibration. Most of our maps are smaller than this, which means, in principle, that some contribution from the

[^17]PSF of the source can basically be found everywhere on the map, and, any emission from the source falling in the background region must be taken into account when estimating the total source flux density.

Taking into account this aperture correction, aperture photometry is performed, using circular apertures of 1.5 times the optical radius whenever possible. For cases where it is not, we adjust our apertures to be sure to encompass all of the FIR emission of the galaxy (Table 4.1). There are three special cases. For HS0052+2536 the chosen aperture also encompasses the neighbouring, very faint galaxy HS0052+2537. Mrk1089 is a galaxy within a compact group of galaxies and UM311 is part of another galaxy and the chosen apertures encompass the whole group of objects. For these galaxies, the spatial resolution of the SPIRE bands makes it very difficult, if not impossible, to separate them from the other objects in their respective groups. For these few cases, the entire group is considered and noted in Tables 3.4, 4.1 and 4.2. The background region is a circular annulus around the source. In most cases, the inner radius of the background region is the same as that of the source aperture and the outer radius is about two times the source aperture radius.

The maps are assumed to consist of the sum of a constant, flat background plus the contribution from the source. Flux densities are measured in the aperture ( $f_{a p}$ ) and in the background annulus $\left(f_{a n}\right)$ by summing the pixels in both regions. The contribution to the measured flux densities ( $f_{a p}$ and $\left.f_{a n}\right)$ from the total flux density of the galaxy $\left(f_{\text {tot }}\right)$ and from the background $(b)$ is determined for each aperture using the encircled energy fraction (eef) tables. These tables, given by HIPE, are measurements of the fraction of the total flux density contained in a given aperture on the PSFs (inverse of the aperture correction). This gives us a simple linear system of two equations with two unknowns: the total flux density from the galaxy $\left(f_{\text {tot }}\right)$ and the background level (b):

$$
\left\{\begin{align*}
f_{a p} & =f_{t o t} \times e e f_{r_{0}}+N_{a p} \times b  \tag{4.1}\\
f_{a n} & =f_{t o t} \times\left(e e f_{r_{2}}-e e f_{r_{1}}\right)+N_{b g} \times b
\end{align*}\right.
$$

where $r_{0}, r_{1}, r_{2}$ are the source aperture radius and the background annulus radii respectively, and $e e f_{r_{0}}, e e f_{r_{1}}$ and $e e f_{r_{2}}$ are the encircled energy fractions at radii $r_{0}, r_{1}, r_{2} . N_{a p}$ (resp. $N_{b g}$ ) is the number of pixels in the source (resp. background) aperture. Inverting this system gives us the values for $f_{t o t}$ and $b$.

If one considers that there is no contribution from the source outside of the source aperture, i.e., setting $e e f_{r_{0}}=1$ and $e e f_{r_{1}}=e e f_{r_{2}}=0$, the flux density will be underestimated, as some contribution from the source will have been removed during the background subtraction. This underestimation depends on the source aperture size, $r_{0}$, and can be important for small apertures. The error made on the flux density becomes greater than the calibration error, which is the dominant source of uncertainty ( $\sim 5 \%$, see Section 4.2.2), when $r_{0} \lesssim 1^{\prime}$. Given that the median $r_{0}$ in the DGS sample is $\sim 0.6^{\prime}$, it is thus important to take into account the contribution from the source falling outside the source aperture.

### 4.2.2 Computing the uncertainties

The uncertainties on the flux density arise from the non-systematic errors due to the measurement of the flux density on the maps, $\left(u n c_{f_{t o t}}\right)$, and the systematic errors due to calibration, ( $\left.u n c_{c a l i b}\right)$.

For the measurement on the maps, the system of equations being linear, the uncertainties arising from the two measurements on the map $\left(u n c_{a p}\right.$ and $\left.u n c_{a n}\right)$ can be linearly propagated to the total flux density and the background level, giving us the uncertainty on the total flux density ( $u n c_{f_{t o t}}$ ) and the uncertainty on the background level $\left(u n c_{b}\right)$. The determination of $u n c_{a p}$ and $u n c_{a n}$ is the same for both errors as the measure is the same: summing pixels in a given region of the map. Thus we detail the calculation for $u n c_{a p}$ only.

There are two sources of errors to $u n c_{a p}$ : one coming from the sum of the pixels, $u n c_{s u m}$, one
coming from the intrinsic error on the flux density value in each pixel, $u n c_{i n t}$.
Determination of $u n c_{s u m}$ : For each pixel there is a contribution from the background noise to the total measured flux density. This error, $\sigma_{s k y}$, is the same for a pixel in the source aperture as well as in the background aperture, repeated $N_{a p}$ times here. The error, $\sigma_{s k y}$, is the standard deviation of all pixels in the background aperture. The final uncertainty, $u n c_{s u m}$, is then:

$$
\begin{equation*}
u n c_{s u m}=\sqrt{N_{a p}} \sigma_{s k y} \tag{4.2}
\end{equation*}
$$

Determination of uncint: For each pixel there is an underlying uncertainty for the flux density value in the pixel, $\sigma_{i n t, i}$, and is independent from pixel to pixel. This uncertainty arises from the data reduction step when the flux density for each pixel is computed. A map of these uncertainties is produced during the data reduction process. The uncertainty, uncint , is then derived by adding quadratically all of the errors in the considered pixels:

$$
\begin{equation*}
u n c_{i n t}=\sqrt{\sum_{i=0}^{N_{a p}} \sigma_{i n t, i}^{2}} \tag{4.3}
\end{equation*}
$$

Note that the assumption of pixel-to-pixel independent uncertainty is not applicable for PACS maps and this can result in an underestimation of $u n c_{i n t}$.

The total error on the source aperture measurement is then:

$$
\begin{equation*}
u n c_{a p}=\sqrt{u n c_{s u m}^{2}+u n c_{i n t}^{2}} \tag{4.4}
\end{equation*}
$$

The $u n c_{a n}$ is derived the same way and we can then compute $u n c_{f_{t o t}}$ and $u n c_{b}$. The quantity $\operatorname{unc}_{f_{\text {tot }}}$ is thus the total error on the flux density due to measurement on the map. To this uncertainty, we add in quadrature the systematic calibration uncertainty, $u n c_{c a l i b}$, of $5 \%$ for the three PACS bands (M. Sauvage \& T. Müller, priv. com.), giving, in the end, the $\sigma_{70-100-160}$ reported in Table 4.1:

$$
\begin{equation*}
\sigma_{\lambda}=\sqrt{u n c_{f_{t o t}}^{2}+u n c_{\text {calib }}^{2}} \tag{4.5}
\end{equation*}
$$

Note that in $u n c_{s u m}$, we have a combination of uncertainties from small-scale astronomical noise and instrumental uncertainties. These instrumental uncertainties can be redundant with part of the instrumental uncertainties taken into account in uncint , leading to an overestimate of $u n c_{a p}$ and thus $u n c_{f_{\text {tot }}}$. However, it has a minor impact on the final uncertainties, $\sigma_{70-100-160}$, as the calibration uncertainty is dominant.

### 4.2.3 Case of upper limits

Some galaxies in our sample are not detected in some or all of the PACS bands. We classify these galaxies as "upper limits" when the computed flux density is lower than five times the corresponding uncertainty on the flux density (e.g., Tol 0618-402, Figure 4.2). We take as the final upper limit, five times the uncertainty on the flux density value in order to have a $5 \sigma$ upper limit (reported in Table 4.1).


Fig. 4.2. Example of a PACS non-detection: (left) PACS $70 \mu \mathrm{~m}$ image of Tol 0618-402. The position of the galaxy is marked with a black cross. The IRAC $8 \mu \mathrm{~m}$ image has been added on the right for comparison. The PACS $70 \mu \mathrm{~m}\left(\mathrm{FWHM}=5.6^{\prime \prime}\right)$ and the IRAC $8 \mu \mathrm{~m}\left(\mathrm{FWHM}=2.0^{\prime \prime}\right)$ beams are indicated as white circles on the bottom right of the images.

### 4.3 DGS SPIRE photometry

For the SPIRE photometre, the RSRFs are different for a point source or for an extended source (see Fig. 3.9 and Section 3.2.2). As explained in Section 3.2.2, the output of the pipeline will be, by definition, a monochromatic flux density of a point source. To obtain monochromatic flux densities of extended sources we apply the ratio of $K_{4}$ corrections for extended and point-like sources, $K_{4 E} / K_{4 P}$, defined in Section 3.2.2. In order to determine which sources need this extra-correction, we have to distinguish between extended and point-like (unresolved) sources in our sample, as well as non-detected sources. Extended sources are defined as galaxies whose spatial extension is larger than the FWHM of the SPIRE beam, and non-detected sources are galaxies that are not visible at SPIRE wavelengths.

### 4.3.1 Extracting the fluxes

The photometry method is adapted for each type of galaxy. However, as the data reduction has been performed with HIPE v5, the $350 \mu \mathrm{~m}$ maps are first scaled by a factor of 1.0067 to update the maps to the latest version of the $350 \mu \mathrm{~m}$ flux calibration (SPIRE Observers' Manual (Section 5.2.8))

## Point source photometry

To determine the flux densities of point sources, we fit a Gaussian function (which is representative of the shape of the PSF) to the timeline data from the bolometers, using a timeline-based source fitter that is used for deriving the flux calibration for the individual bolometers ${ }^{4}$. We then check a posteriori that our "unresolved" classification was correct: if the FWHM of the fitted Gaussian is $<20^{\prime \prime}, 29^{\prime \prime}$ and $37^{\prime \prime}$ at 250,350 and $500 \mu \mathrm{~m}$ respectively, then the source can be considered as truly point-like. As the timeline data is in Jy beam ${ }^{-1}$, the flux density will simply be the amplitude of the fitted Gaussian. This is the most accurate way of computing flux densities for point-like sources as it matches the measurement techniques used for the SPIRE calibration. Moreover we avoid all pixelization issues when using the timeline data rather than the map. On top of that, applying any

[^18]map-making process would also smear the PSFs, causing the peak signal values to decrease by $\sim$ $5 \%$ for point sources.

## Extended source photometry

For the extended sources, we perform aperture photometry on the maps, using the same source and background apertures as those used for the PACS photometry, and check that the PACS apertures do fully encompass the SPIRE emission from the entire galaxy. The maps are converted from Jy beam $^{-1}$ to Jy pix ${ }^{-1}$ considering that the beam area values are 465,822 and 1768 square arcseconds ${ }^{5}$ at $250,350,500 \mu \mathrm{~m}$ respectively and the pixel sizes are given in Section 4.1.2.

The background level is determined by the median of all of the pixels in the background aperture. The median is preferred rather than the mean because the SPIRE background is contaminated by prolific background sources due to some observations reaching the confusion limit. The background level is then subtracted from our maps and the total flux density is the sum of all of the pixels encompassed in the source aperture, corrected for $K_{4 E} / K_{4 P}$. These $K_{4 E} / K_{4 P}$ correction factors, given in the SPIRE Observers' Manual (Section 5.2.8), are 0.98279, 0.98344 and 0.97099 at 250, 350, $500 \mu \mathrm{~m}$ respectively.

However there are also "marginally" extended sources (e.g., IIZw40) that do not require this $K_{4 E} / K_{4 P}$ correction. To identify these sources, we first check that the source is truly resolved by applying the point source method on the timeline data. We verify that the FWHM is indeed greater than the chosen threshold values for the "unresolved" classification. As an additional check, the fitted Gaussian is subtracted from the map and the resulting map is visually checked for any remaining emission from the source. If this condition is satisfied, then the source is truly resolved. If the FWHM of the fitted Gaussian is lower than $24^{\prime \prime}, 34^{\prime \prime}$ and $45^{\prime \prime}$ at $250,350,500 \mu \mathrm{~m}$ respectively then the source is considered to be "marginally" extended only, and thus to not require the $K_{4 E} / K_{4 P}$ correction.

### 4.3.2 Computing the uncertainties

As for the PACS photometry, there are two types of uncertainties for SPIRE photometry: the errors arising from the determination of the flux density, unc $c_{f l u x}$, and the calibration errors, $u n c_{c a l i b}$. As we used different methods for flux extraction depending on the type of the source, the errors contributing to $u n c_{f l u x}$ are determined differently. The method described here has been adapted from the method described in Ciesla et al. (2012).

## Point sources

The uncertainty on the flux density for a point source is determined through a test in which we add 100 artificial sources with the same flux density as the original source. They are added at random locations in the maps, within a 0.3 deg box centred on the original source. The same photometry procedure was applied to the artificial sources and the final uncertainty is the standard deviation in the flux densities derived for the artificial sources. We quote the following uncertainties (unc flux $)$ for point-like sources:

- 6 mJy at $250 \mu \mathrm{~m}$;
- 7 mJy (for flux densities $>50 \mathrm{mJy}$ ) and 10 mJy (for flux densities $\lesssim 50 \mathrm{mJy}$ ) at $350 \mu \mathrm{~m}$;
- 9 mJy at $500 \mu \mathrm{~m}$.

[^19]
## Extended sources

For the aperture photometry performed on the extended sources, we have four types of uncertainties contributing to $u n c_{\text {flux }}$ :

- the uncertainty arising from the background level determination, $u n c_{b g}$,
- the uncertainty due to background noise in the source aperture, $u n c_{\text {source }}$,
- the underlying uncertainty for the flux density value in the pixel coming from the data reduction, $u n c_{i n t}$, and
- the uncertainty in the beam area value: $u n c_{\text {beam }}$, which is given to be $4 \%^{6}$.

The determination of the background level generates an uncertainty that will affect each pixel in the source aperture when subtracting the background level from the map. The uncertainty on the background level is $u n c_{b g_{l e v e l}}=\sigma_{s k y} / \sqrt{N_{b g}}$, with $\sigma_{s k y}$ being here again the standard deviation of all of the pixels in the background aperture. This will affect the determination of the flux density for each pixel summed in the aperture:

$$
\begin{equation*}
u n c_{b g}=N_{a p} u n c_{b g_{l e v e l}} \tag{4.6}
\end{equation*}
$$

The uncertainty due to background noise in the source aperture, $u n c_{\text {source }}$, is determined the same way as the PACS $u n c_{a p}$ since it is the uncertainty arising from summing the pixels in a given aperture:

$$
\begin{equation*}
u n c_{s o u r c e}=\sqrt{N_{a p}} \sigma_{s k y} \tag{4.7}
\end{equation*}
$$

The uncertainty arising from the underlying uncertainties of the flux density value in each pixel is computed the same way as for PACS. Here again, this uncertainty arises from the data reduction step when the flux density for each pixel is computed, and the pipeline produces the corresponding error map:

$$
\begin{equation*}
u n c_{i n t}=\sqrt{\sum_{i=0}^{N_{a p}} \sigma_{i n t, i}^{2}} \tag{4.8}
\end{equation*}
$$

The total uncertainty coming from the determination of the flux density for an extended source, is then:

$$
\begin{equation*}
u n c_{f l u x}=\sqrt{u n c_{b g}^{2}+u n c_{\text {source }}^{2}+u n c_{i n t}^{2}+u n c_{\text {beam }}^{2}} \tag{4.9}
\end{equation*}
$$

For both types of sources, we also add calibration uncertainties to $u n c_{f l u x}$ to get the final total uncertainty. There are two different SPIRE calibration uncertainties: a systematic uncertainty of $\sim$ $5 \%$ coming from the models used for Neptune, the primary calibrator, which is correlated between the three bands, and a random uncertainty of $\sim 2 \%$ coming from the repetitive measurement of the flux densities of Neptune. These two uncertainties were not added in quadrature but linearly, as advised in the SPIRE Observer's Manual, giving an overall $7 \%$ calibration uncertainty unc calib. The final total uncertainty, $\sigma_{250-350-500}$ reported in Table 4.2 , is obtained by adding $u n c_{f l u x}$ and $u n c_{c a l i b}$ in quadrature.

As for PACS, with SPIRE we also have a redundancy in the error estimation in unc source and $u n c_{i n t}$, again with only a minor impact on the final uncertainties, $\sigma_{250-350-500}$, as the calibration uncertainty dominates.

[^20]

Fig. 4.3. Example of a SPIRE non-detection: (left) SPIRE $250 \mu \mathrm{~m}$ and (right) PACS $70 \mu \mathrm{~m}$ image of SBS 0335-052. The position of the galaxy is indicated by a black cross on the SPIRE image. The SPIRE $250 \mu \mathrm{~m}$ ( $\mathrm{FWHM}=18.2^{\prime \prime}$ ) and the PACS $70 \mu \mathrm{~m}\left(\mathrm{FWHM}=5.6^{\prime \prime}\right)$ beams are indicated as white circles on the bottom right of the images.

### 4.3.3 Case of upper limits

When the galaxy is not detected in the SPIRE bands (e.g., SBS 0335-052, Figure 4.3), we can only derive upper limits on the flux density. Also, when the source is blended with another source in the beam and we are unable to confidently separate them (e.g., Pox186 and a background galaxy separated by $20^{\prime \prime}$, Figure 4.4), upper limits are reported. Since the undetected sources are point sources, we use five times the uncertainties reported for point sources in Section 4.3.2. The only exception is SBS1533+574 which is blended with another source and slightly extended at $250 \mu \mathrm{~m}$. The method described above gives an upper limit too low. The extended source photometry method is thus used to derive a $5 \sigma$ upper limit at this wavelength.

### 4.3.4 Special cases: heavy cirrus contamination

For NGC 6822 and IC 10, the cirrus contamination from our Galaxy is significant in the SPIRE bands.

NGC 6822 - Galametz et al. (2010) determined that the contribution from the cirrus to the total emission of the galaxy is of the order of $30 \%$ for all SPIRE bands. To determine the cirrus contribution here, we assume that the entire galaxy is in a homogeneous and flat cirrus region. We determine this cirrus level by considering regions at the same cirrus level outside of the galaxy. This level is used as the background level for the flux determination. We then compare this flux density with the flux density obtained when we consider an uncontaminated background region and determine the contamination from the cirrus. We also find that the contribution of the cirrus to the total flux densities is about $30 \%$, which is coherent with the results from Galametz et al. (2010). Thus for this galaxy, the flux densities cited in Table 4.2 are flux densities where the cirrus contribution has been subtracted. We also include a conservative $30 \%$ uncertainty in the error for these flux densities to account for the estimation of the cirrus contribution, and for the fact that the cirrus emission may not be flat.

IC 10 - We apply the same method here. Again, we find that the cirrus contributes $\sim 30 \%$ on average, to each SPIRE band. We took this contribution into account by adding this cirrus


Fig. 4.4. Example of a "mixed" source. SPIRE $500 \mu \mathrm{~m}$ (left) and PACS 160 (right) images of Pox186 and a contaminating background source. The sources are $20^{\prime \prime}$ apart, and are well separated at $160 \mu \mathrm{~m}$, but are completely blended at SPIRE $500 \mu \mathrm{~m}$ resolution. Pox 186 corresponds to the bottom cross, whereas the contaminating background source is the X. The SPIRE $500 \mu \mathrm{~m}$ (FWHM=36.3") and the PACS $160 \mu \mathrm{~m}$ ( $\mathrm{FWHM}=11.3^{\prime \prime}$ ) beams are indicated as white circles on the bottom right of the images.
uncertainty to the other sources of uncertainties for this galaxy.
This method can be improved, by using the HI maps to better determine the cirrus emission and the background level and thus reducing the uncertainties on the measurements for these two galaxies.

### 4.4 KINGFISH data

The Herschel KINGFISH flux densities are taken from Dale et al. (2012). They have been extracted from scanamorphos maps. The KINGFISH SPIRE fluxes and corresponding uncertainties are updated to match the latest SPIRE beam areas (see Section 4.3.1). The beam areas used in this paper were released in September 2012, after the publication of Dale et al. (2012) in January 2012.
Table 4.1. Table of PACS flux densities for the DGS galaxies. The map-making method is indicated for each galaxy as well as the radius of the circular aperture used for the flux extraction. When an upper limit is given, it is the $5 \sigma$ upper limit computed in 4.2.3.

| Source | $\mathrm{F}_{70}$ (Jy) | $\sigma_{70}$ (Jy) | $\mathrm{F}_{100}$ (Jy) | $\sigma_{100}$ (Jy) | $\mathrm{F}_{160}$ (Jy) | $\sigma_{160}$ (Jy) | Method | Circular aperture radius ( ${ }^{\prime \prime}$ ) |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Harol1 | 6.14 | 0.31 | 4.96 | 0.25 | 2.42 | 0.12 | SCANAMORPHOS | $45^{6}$ |
| Haro2 | 4.99 | 0.25 | 5.33 | 0.27 | 3.95 | 0.20 | Scanamorphos | $50^{a}$ |
| Haro3 | 5.30 | 0.26 | 6.41 | 0.32 | 4.83 | 0.24 | Scanamorphos | $60^{a}$ |
| He2-10 | 25.6 | 1.3 | 26.6 | 1.3 | 18.8 | 0.9 | Scanamorphos | $108^{a}$ |
| HS0017+1055 | 0.046 | 0.005 | 0.033 | 0.004 | 0.019 | 0.004 | РhotProject | $16^{6}$ |
| HS0052+2536 | 0.22 | 0.01 | 0.21 | 0.01 | 0.139 | 0.008 | РhotProject | $23^{a, f}$ |
| HS0822+3542 | $\leq 0.014^{j} / \leq 0.041^{k}$ | - | $\leq 0.013^{j} / \leq 0.048^{k}$ | - | 0.034 | 0.003 | PhotProject | $12^{a}$ |
| HS1222+3741 | 0.025 | 0.004 | $\leq 0.036$ | - | $\leq 0.022$ | - | РhotProject | $14^{\text {b }}$ |
| HS1236+3937 | $\leq 0.029$ | - | $\leq 0.035$ | - | $\leq 0.028$ | - | РhotProject | $15^{a}$ |
| HS1304+3529 | 0.121 | 0.007 | ${ }^{-150}$ | 0.009 | ${ }^{-} 0.069$ | 0.005 | РhotProject | $18^{6}$ |
| HS1319+3224 | 0.012 | 0.003 | 0.013 | 0.002 | $\leq 0.015$ | - | РhotProject | $8^{\text {b }}$ |
| HS1330+3651 | 0.093 | 0.006 | 0.112 | 0.007 | 0.091 | 0.005 | РhotProject | $20^{\text {b }}$ |
| HS1442+4250 | 0.09 | 0.01 | $\leq 0.016$ | - | $\leq 0.047$ | - | РhotProject | $51^{a}$ |
| HS2352+2733 | 0.039 | 0.003 | 0.016 | 0.002 | $\leq 0.016$ | - | РhotProject | $15^{a}$ |
| IZw18 | 0.045 | 0.003 | 0.018 | 0.002 | $\leq 0.011$ | - | РhotProject | $14^{a}$ |
| IC10 | 140. | 7. | 207. | 10. | 225. | 11. | Scanamorphos | $306{ }^{\text {a }}$ |
| IIZw40 | 6.39 | 0.32 | 5.79 | 0.29 | 3.53 | 0.18 | Scanamorphos | $66^{b}$ |
| Mrk1089 | 4.27 | 0.21 | 4.97 | 0.25 | 4.68 | 0.23 | Scanamorphos | $75^{\text {b,f }}$ |
| Mrk1450 | 0.30 | 0.02 | 0.25 | 0.01 | 0.127 | 0.007 | РнотProject | $20^{a}$ |
| Mrk153 | 0.28 | 0.02 | 0.30 | 0.02 | 0.137 | 0.009 | РhotProject | $35^{b}$ |
| Mrk209 | 0.32 | 0.02 | 0.35 | 0.02 | 0.16 | 0.01 | Scanamorphos | $24^{c}$ |
| Mrk930 | 1.19 | 0.06 | 1.40 | 0.07 | 0.98 | 0.05 | Scanamorphos | $60^{\text {b }}$ |
| NGC1140 | 4.04 | 0.20 | 4.62 | 0.23 | 4.58 | 0.23 | Scanamorphos | $118^{\text {b }}$ |
| NGC1569 | 60.4 | 3.0 | 57.3 | 2.9 | 39.7 | 2.0 | SCANAMORPHOS | $150{ }^{\text {d }}$ |
| NGC1705 | 1.37 | 0.07 | 1.46 | 0.07 | 1.10 | 0.06 | РhotProject ${ }^{g}$ | $72^{d}$ |
| NGC2366 | 5.30 | 0.26 | 6.23 | 0.31 | 4.08 | 0.20 | Scanamorphos | $150^{\text {d,e }}$ |
| NGC4214 | 24.5 | 1.2 | 32.2 | 1.6 | 33.7 | 1.7 | Scanamorphos | $300{ }^{\text {d }}$ |
| NGC4449 | 49.3 | 2.5 | 75.9 | 3.8 | 79.5 | 4.0 | Scanamorphos | $250{ }^{\text {d }}$ |
| NGC4861 | 2.31 | 0.12 | 2.17 | 0.11 | 1.99 | 0.10 | SCANAMORPHOS | $120^{\text {d }}$ |
| NGC5253 | 32.9 | 1.6 | 32.3 | 1.6 | 23.2 | 1.2 | Scanamorphos | $120^{\text {d }}$ |
| NGC625 | 6.49 | 0.32 | 9.47 | 0.47 | 8.20 | 0.41 | Scanamorphos | $170^{\text {d }}$ |
| NGC6822 | 54.9 | 2.8 | 63.6 | 3.2 | 77.1 | 3.9 | Scanamorphos | $440^{\text {d }}$ |
| Pox186 | $0.038^{j}$ | 0.005 | $0.052^{j}$ | 0.005 | $0.047^{j}$ | 0.004 | РhotProject | $16^{a}$ |
| SBS0335-052 | 0.056 | 0.004 | $0.024^{h}$ | 0.001 | $0.007^{h}$ | 0.001 | РhotProject | $10^{a, h}$ |
| SBS1159+545 | 0.019 | 0.003 | 0.019 | 0.003 | $\leq 0.018$ | - | РhotProject | $8^{a}$ |
| SBS1211+540 | 0.034 | 0.003 | 0.018 | 0.002 | 0.013 | 0.002 | РhotProject | $15^{\text {b }}$ |
| SBS1249+493 | 0.032 | 0.005 | $\leq 0.034$ | - | $\leq 0.042$ | - | РhotProject | $12^{\text {b }}$ |
| SBS1415+437 | 0.18 | 0.01 | 0.16 | 0.01 | 0.065 | 0.007 | РhotProject | $34^{a}$ |
| SBS1533+574 | 0.19 | 0.01 | 0.24 | 0.01 | 0.19 | 0.01 | Scanamorphos | $30^{a}$ |
| Tol0618-402 | $\leq 0.014$ | - | $\leq 0.005$ | - | $\leq 0.013$ | - | РhotProject | $18^{a}$ |
| Tol1214-277 | 0.017 | 0.003 | 0.018 | 0.002 | $\leq 0.018$ | - | РhotProject | $12^{\text {b }}$ |
| UGC4483 | 0.16 | 0.02 | ${ }_{-i}$ | $-{ }^{i}$ | $\leq 0.037^{j} / \leq 0.072^{k}$ | - | РhotProject | $63^{a}$ |
| UGCA20 | $\leq 0.052$ | . | $\leq 0.057$ | - | $\leq 0.048$ | - | РhotProject | $20^{\text {c }}$ |

Table 4.1. Table of PACS flux densities for the DGS galaxies (continued).

| Source | $\mathrm{F}_{70}(\mathrm{Jy})$ | $\sigma_{70}(\mathrm{Jy})$ | $\mathrm{F}_{100}(\mathrm{Jy})$ | $\sigma_{100}(\mathrm{Jy})$ | $\mathrm{F}_{160}(\mathrm{Jy})$ | $\sigma_{160}(\mathrm{Jy})$ | Method | Circular aperture radius ( $\left.{ }^{\prime \prime}\right)$ |
| :--- | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| UM133 | $0.155^{j}$ | 0.01 | 0.066 | 0.010 | 0.053 | 0.009 | SCANAMORPHOS | $26^{c}$ |
| UM311 | 2.94 | 0.15 | 5.63 | 0.28 | 6.10 | 0.31 | SCANAMORPHOS | $140^{b, f}$ |
| UM448 | 5.17 | 0.26 | - | - | 3.22 | 0.17 | SCANAMORPHOS | $64^{b}$ |
| UM461 | $0.21^{j}$ | 0.01 | 0.145 | 0.009 | 0.113 | 0.007 | PhotProJect | $17^{b}$ |
| VIIZw403 | 0.47 | 0.03 | 0.56 | 0.03 | 0.34 | 0.02 | PhotProJect | $40^{c}$ |

${ }^{a, b, c}$ : The radius is: ${ }^{a}$ : 1.5 times the optical radius. ${ }^{b}$ : larger than 1.5 times the optical radius. ${ }^{c}$ : smaller than 1.5 times the optical radius.
$f$ : These objects are galaxies within compact groups of galaxies or are parts of other galaxies and the photometry given here is for the whole group (see Section 4.2 .1 for details). ${ }^{g}$ : The Scanamorphos maps of NGC1705 are not satisfactory because of a non uniform background. Therefore we use the PhotProject maps. To preserve as much of the diffuse
extended emission as possible we manually mask out the galaxy before performing the high-pass filtering step.
: New observations were obtained for SBS $0335-052$ at 100 and $160 \mu \mathrm{~m}$ with longer integration times (Sauvage et al. in prep). We chose to quote the flux density values from the newest observations for 100 and $160 \mu \mathrm{~m}$.
$i$
: Interferences on the detector are stron
galaxy.
$j$ : These flux densities might present some discrepancies with other FIR measurements (i.e., MIPS, other PACS and/or SPIRE wavelengths).
${ }^{k}$ : These fluxes densities are more coherent with the rest of the IR photometry for this galaxy (see Section 7.2.4). Thus we recommend to use these values when studying this object.

Table 4.2. Table of SPIRE flux densities for the DGS galaxies. When an upper limit is given, it is the $5 \sigma$ upper limit computed in 4.3.3.

| Source | $\mathrm{F}_{250}$ (Jy) | $\sigma_{250}$ (Jy) | $\mathrm{F}_{350}$ (Jy) | $\sigma_{350}(\mathrm{Jy})$ | $\mathrm{F}_{500}$ (Jy) | $\sigma_{500}(\mathrm{Jy})$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Haro11 | 0.63 | 0.05 | 0.23 | 0.02 | 0.09 | 0.01 |
| Haro2 | $1.28{ }^{\text {a }}$ | 0.10 | $0.53{ }^{\text {a }}$ | 0.04 | $0.15{ }^{\text {a }}$ | 0.01 |
| Haro3 | $1.79{ }^{\text {a }}$ | 0.15 | $0.77^{a}$ | 0.07 | $0.23{ }^{\text {a }}$ | 0.02 |
| He2-10 | $6.67{ }^{a}$ | 0.54 | $2.64{ }^{a}$ | 0.22 | $0.79^{a}$ | 0.07 |
| HS0017+1055 | $\leq 0.030$ | - | $\leq 0.050$ | - | $\leq 0.045$ | - |
| HS0052+2536 ${ }^{\text {b }}$ | 0.058 | 0.007 | 0.03 | 0.01 | 0.018 | 0.009 |
| HS0822+3542 ${ }^{\text {c }}$ | - | - | - | - | - | - |
| HS1222+3741 ${ }^{\text {c }}$ | - | - | - | - | - | - |
| HS1236+3937 | $\leq 0.030$ | - | $\leq 0.050$ | - | $\leq 0.045$ | - |
| HS1304+3529 | 0.038 | 0.007 | $\leq 0.050$ | - | $\leq 0.045$ | - |
| HS1319+3224 ${ }^{\text {c }}$ | - | - | - | - | - | - |
| HS1330+3651 ${ }^{\text {c }}$ | - | - | - | - | - | - |
| HS1442+4250 | $\leq 0.030$ | - | $\leq 0.050$ | - | $\leq 0.045$ | - |
| HS2352+2733 | $\leq 0.030$ | - | $\leq 0.050$ | - | $\leq 0.045$ | - |
| IZw18 | $\leq 0.030$ | - | $\leq 0.050$ | - | $\leq 0.045$ | - |
| IC10 ${ }^{\text {d }}$ | $101 .{ }^{\text {a }}$ | 31. | $47.6{ }^{\text {a }}$ | 14.8 | $16.3{ }^{a}$ | 5.1 |
| IIZw40 | $1.33{ }^{a}$ | 0.12 | $0.58{ }^{a}$ | 0.06 | 0.18 | 0.02 |
| Mrk1089 ${ }^{\text {b }}$ | $1.75{ }^{\text {a }}$ | 0.15 | $0.78{ }^{a}$ | 0.07 | $0.24{ }^{a}$ | 0.03 |
| Mrk1450 | 0.049 | 0.007 | $\leq 0.050$ | - | $\leq 0.045$ | - |
| Mrk153 | $0.048^{a}$ | 0.008 | $\leq 0.050$ | - | $\leq 0.045$ | - |
| Mrk209 | 0.062 | 0.007 | 0.03 | 0.01 | $\leq 0.045$ | - |
| Mrk930 | $0.40{ }^{\text {a }}$ | 0.04 | $0.13{ }^{a}$ | 0.01 | $0.049^{a}$ | 0.007 |
| NGC1140 | $1.97{ }^{\text {a }}$ | 0.17 | $0.94{ }^{a}$ | 0.08 | $0.28^{a}$ | 0.03 |
| NGC1569 | $12.0^{a}$ | 1.0 | $5.02{ }^{\text {a }}$ | 0.41 | $1.55{ }^{\text {a }}$ | 0.13 |
| NGC1705 | $0.60{ }^{\text {a }}$ | 0.05 | $0.29{ }^{\text {a }}$ | 0.03 | $0.10^{a}$ | 0.01 |
| NGC2366 | $2.04{ }^{\text {a }}$ | 0.17 | $1.01{ }^{\text {a }}$ | 0.09 | $0.39^{a}$ | 0.04 |
| NGC4214 | $18.6^{a}$ | 1.5 | $9.92{ }^{a}$ | 0.80 | $3.79^{a}$ | 0.31 |
| NGC4449 | $32.4{ }^{\text {a }}$ | 2.6 | $14.8{ }^{\text {a }}$ | 1.2 | $5.01{ }^{\text {a }}$ | 0.41 |
| NGC4861 | $1.10^{a}$ | 0.10 | $0.54{ }^{a}$ | 0.05 | $0.20^{a}$ | 0.03 |
| NGC5253 | $7.82{ }^{\text {a }}$ | 0.63 | $3.64{ }^{\text {a }}$ | 0.29 | $1.18{ }^{a}$ | 0.10 |
| NGC625 | $4.33{ }^{\text {a }}$ | 0.35 | $2.18^{a}$ | 0.18 | $0.80^{a}$ | 0.07 |
| NGC6822 ${ }^{\text {d }}$ | $48.4{ }^{\text {a }}$ | 15.0 | $29.7^{a}$ | 9.2 | $13.6{ }^{a}$ | 4.2 |
| Pox186 | 0.045 | 0.007 | $\leq 0.050$ | - | $\leq 0.045$ | - |
| SBS0335-052 | $\leq 0.030$ | - | $\leq 0.050$ | - | $\leq 0.045$ | - |
| SBS1159+545 | $\leq 0.030$ | - | $\leq 0.050$ | - | $\leq 0.045$ | - |
| SBS1211+540 | $\leq 0.030$ | - | $\leq 0.050$ | - | $\leq 0.045$ | - |
| SBS1249+493 | $\leq 0.030$ | - | $\leq 0.050$ | - | $\leq 0.045$ | - |
| SBS1415+437 ${ }^{\text {c }}$ | - | - | - | - | - | - |
| SBS1533+574 | $\leq 0.122^{a}$ | - | $\leq 0.050$ | - | $\leq 0.045$ | - |
| Tol0618-402 ${ }^{\text {c }}$ | - | - | - | - | - | - |
| Tol1214-277 | $\leq 0.030$ | - | $\leq 0.050$ | - | $\leq 0.045$ | - |
| UGC4483 | 0.024 | 0.006 | $\leq 0.050$ | - | $\leq 0.045$ | - |
| UGCA20 ${ }^{\text {c }}$ | - | - | - | - | - | - |
| UM133 | 0.032 | 0.006 | $\leq 0.050$ | - | $\leq 0.045$ | - |
| UM311 ${ }^{\text {b }}$ | $3.84{ }^{a}$ | 0.31 | $1.87^{a}$ | 0.16 | $0.66{ }^{a}$ | 0.06 |
| UM448 | $0.99^{a}$ | 0.08 | $0.38{ }^{\text {a }}$ | 0.03 | 0.13 | 0.01 |
| UM461 | 0.027 | 0.006 | 0.03 | 0.01 | $\leq 0.045$ | - |
| VIIZw403 | $0.14{ }^{a}$ | 0.01 | 0.053 | 0.008 | 0.028 | 0.009 |

${ }^{a}$ : The flux densities are derived from aperture photometry, with the same aperture used for PACS.
${ }^{b}$ : These objects are galaxies within compact groups of galaxies or are parts of other galaxies and the photometry given here is for the whole group (see Section 4.2 .1 for details).
${ }^{c}$ : These sources were not observed at all by SPIRE.
${ }^{d}$ : The quoted flux densities for these sources have been corrected for cirrus contamination.

## Chapter 5

## Cold dust from low-metallicity environments to normal galaxies

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This Chapter is dedicated to the comparison the cold dust properties of dwarf galaxies with that of more metal-rich environments, first qualitatively with colour-colour diagrams (Section 5.1), and then quantitatively with the aid of a modified blackbody model (Section 5.2). The cold dust properties are thus characterised with the following parameters: the temperature ( $T$ ), the emissivity index $(\beta)$, the dust mass ( $M_{d u s t, B B}$ ) and the FIR luminosity ( $L_{F I R}$ ).

### 5.1 Characterising the SED shapes

In order to obtain a qualitative view of the FIR-to-submm behaviour of the DGS sample, and to compare with the KINGFISH sample, we inspect the observed Herschel SEDs as well as several Herschel colour-colour diagrams combining both PACS and SPIRE observations. Indeed, recent studies have shown how colours can be used to explore the dust properties of galaxies (e.g., Bendo et al. 2010a; Galametz et al. 2010; Boselli et al. 2010a, 2012; Dale et al. 2012; Galametz et al. 2013).

### 5.1.1 Observed spectral energy distributions



Fig. 5.1. Total Herschel observed SEDs for both DGS and KINGFISH samples, normalized at $70 \mu \mathrm{~m}$. The colours delineate the different metallicity bins, and the lines and symbols differentiate DGS (plain lines and crosses) and KINGFISH galaxies (dashed lines and downward triangles).

Total observed SEDs for both samples are computed for a first look at the characteristic SED shapes in the DGS and KINGFISH samples (Figure 5.1). The upper limits are not indicated here for clarity. The most metal-poor galaxies are also the faintest and therefore not detected with

Herschel beyond $160 \mu \mathrm{~m}$. The observed SEDs are normalized at $70 \mu \mathrm{~m}$, and we see here that the peak of the SED shifts towards longer wavelengths as the metallicity increases, reflecting the impact of metallicity on the observed dust properties.

### 5.1.2 Dwarf Galaxy Survey colours

## Constructing the colour-colour diagrams

The Herschel colour-colour diagrams are constructed first by computing the observed flux ratios and the corresponding error bars, for both DGS and KINGFISH, and omitting the galaxies with more than one upper limit in the considered bands.

We then compute the theoretical Herschel flux ratios of simulated modified blackbodies spanning a range in temperature ( $T$ from 0 to 40 K in 2 K bins and from 40 to 100 K in 10 K bins) and emissivity indices ( $\beta$ from 0.0 to 2.5 ). From now on, we define the emissivity index fixed for modelling the simulated Herschel flux ratios as " $\beta_{\text {theo }}$ ", and " $\beta_{\text {obs }}$ " when we leave the emissivity index as a free parameter in modified blackbody fits (see Section 5.2). In our simulated modified blackbody, the emitted fluxes are proportional to $\lambda^{-\beta_{\text {theo }}} \times B_{\nu}(\lambda, T)$, where $B_{\nu}(\lambda, T)$ is the Planck function (see Section 5.2).

The pipeline we use for the data reduction gives us monochromatic flux densities for our data points for both PACS and SPIRE. To mimic the output of the pipeline for our theoretical points we weight our theoretical flux density estimates by the RSRF of the corresponding bands. For SPIRE simulated measurements, we then convert our RSRF-weighted flux densities into monochromatic flux densities by applying the $K_{4}$ correction given on the SPIRE Observers' Manual (see Section 3.2.2). For PACS, we also colour correct the RSRF-weighted modelled flux densities to a spectrum where $\nu \mathrm{F}_{\nu}$ is constant (i.e., multiply by the analogous of $K_{4}$ for PACS, see Eq. 3.4). These simulated flux ratios from a simple model are useful indicators to interpret the colour-colour diagrams.

## FIR/submm colours

The spread of galaxies on the colour-colour diagrams (Figures 5.2 and 5.3) reflects broad variations in the SED shape and physical properties in our survey.

The $F_{70} / F_{100}$ vs $F_{100} / F_{160}$ diagram (Figure 5.2) traces best the peak of the SED. Galaxies usually exhibit a peak in their SED around $\sim 100-160 \mu \mathrm{~m}$. Galaxies with FIR flux densities $F_{70}>F_{100}>F_{160}$ may be quite warm as they peak at wavelengths less than $70 \mu \mathrm{~m}$. Colder galaxies would lie in the lower left corner of the plot ( $F_{70}<F_{100}<F_{160}$ ), as shown by the simulated flux ratio lines. KINGFISH galaxies indeed cluster in the corresponding lower-left corner of the plot while DGS galaxies span a wider space (Fig 5.2, top), reflecting the differences in the dust properties between the low-metallicity dwarf galaxies and the generally more metal-rich environments probed by the KINGFISH survey. Nonetheless both samples follow the trends outlined by the theoretical flux ratio lines from simulated modified blackbodies. There are some outliers, all of them being very faint, extremely metal-poor galaxies (from 0.03 to $0.20 Z_{\odot}$ ). There is also a metallicity trend in Fig. 5.2 (bottom), between the KINGFISH sample and the DGS galaxies, as well as within both samples, i.e., low-metallicity (dwarf) galaxies peak at much shorter wavelengths and thus have overall warmer dust (several tens of K ), compared to more metal-rich galaxies.

As we saw in Section 2.2.2, the warmer dust of dwarf galaxies is due to the very energetic environment in which the grains reside: the dust grains are exposed to harder and more intense ISRF than in a more metal-rich environment. This increases the contribution of hot and warm dust to the total dust emission resulting in higher equilibrium dust temperatures. This confirms what has already been observed by $\operatorname{IRAS}$ and Spitzer on smaller samples of dwarf galaxies (e.g., Hunter et al. 1989; Sauvage et al. 1990; Melisse \& Israel 1994; Walter et al. 2007).


Fig. 5.2. Colour-colour diagram: PACS/PACS diagram: $F_{70} / F_{100}$ versus $F_{100} / F_{160}$. (top) The colour and symbol code differentiates DGS (purple crosses) and KINGFISH galaxies (orange downward triangles). (bottom) The colour code delineates the different metallicity bins this time. Crosses and downward triangles are still representing DGS and KINGFISH galaxies, respectively. For both plots, the curves give theoretical Herschel flux ratios for simulated modified blackbodies for $\beta_{\text {theo }}=0.0$ to 2.5 and $T$ from 0 to 40 K in 2 K bins and from 40 to 100 K in 10 K bins, as black dots, increasing in $T$ from left to right. Lines of constant $T$ are indicated as dotted lines, and a few temperatures have been marked on the plots.


Fig. 5.3. Colour-colour diagram: PACS/SPIRE diagram: $F_{100} / F_{250}$ versus $F_{250} / F_{500}$. The colour and symbol choices are the same as in Figure 5.2 for both figures. Note that the most metal-poor galaxies (from 0.03 to $\left.0.20 Z_{\odot}\right)$ are very faint and even not detected anymore at long wavelengths. We were only able to derive upper limits beyond $160 \mu \mathrm{~m}$ for these galaxies and, thus, some galaxies do not appear on this diagram anymore.

Note that there is a small excess at $70 \mu \mathrm{~m}$ for most of the galaxies compared to our simulated modified blackbodies, causing them to fall above the lowest $\beta_{\text {theo }}$ line to artifically low $\beta_{\text {obs }}$. This means that if we were to fit modified blackbodies only to the FIR flux densities (from $70 \mu \mathrm{~m}$ to 160 $\mu \mathrm{m}$ ) we would get very low values of $\beta_{o b s}$, i.e., very flat SEDs in the FIR, which reflects a broad peak in the observed SED. This is due to the crudeness of the isothermal approximation made in the modified blackbody modelling. In a real galaxy, the dust grains have a range of temperatures, (e.g., hotter dust around star-forming regions vs colder dust in the diffuse ISM). Such a low $\beta_{o b s}$ here is only a side effect of the distribution in temperature of the grains in the galaxy. The extremely metalpoor outliers noted before may have an even wider temperature distribution than in more metal-rich galaxies, causing the broadening of the peak in their dust SED and their peculiar location on the diagrams in Fig. 5.2. Part of this excess at $70 \mu \mathrm{~m}$ could also be due to non-thermal heating, i.e., dust grains whose emission can not be represented by a modified blackbody, such as stochastically heated small grains.

More accurate values of $T$ and $\beta_{\text {theo }}$ may be illustrated by including submm data in the colourcolour diagrams. At submm wavelengths (beyond $\sim 250 \mu \mathrm{~m}$ ), the emissivity index this time represents an intrinsic grain property: the efficiency of the emission from the dust grain. A theoretical emissivity index $\beta_{\text {theo }}=2$ is commonly used to describe the submm SED for local and distant galaxies in the models as it represents the intrinsic optical properties of Galactic grains (mixture of graphite and silicate grains; see Sections 1.2 .3 and 1.2.5). More recently $\beta_{\text {theo }}$ between 1.5 and 2 have also been used for galaxies (e.g., Amblard et al. 2010; Dunne et al. 2011). The $F_{100} / F_{250}$ vs $F_{250} / F_{500}$ diagram (Figure 5.3) reflects best the variations in emissivity index $\beta_{\text {theo }}$. Here again the DGS galaxies are more widely spread than the KINGFISH galaxies (Figure 5.3, top) spanning larger ranges of $F_{100} / F_{250}$ and $F_{250} / F_{500}$ ratios, that is, wider ranges in temperature and $\beta$ (such as higher $T$ and lower $\beta$ ). As far as metallicity is concerned, the trend with temperature already noted in Fig. 5.2 is still present (Figure 5.3, bottom). But hardly any trend between $\beta$ and metallicity can be noticed: as the extremely low-metallicity galaxies are not detected at $500 \mu \mathrm{~m}$, it is rather difficult to conclude on this point relying only on the FIR/submm colour-colour diagram.

Modelling low-metallicity dwarf galaxies with grain properties derived from the Galaxy (i.e., using $\beta_{\text {theo }}=2$ ), may thus not be appropriate. The galaxies showing a lower $\beta_{o b s}\left(\beta_{o b s}<2\right)$ will have a flatter submm slope. Smaller $F_{250} / F_{500}$ ratios, that can be seen as a sign of lower $\beta_{\text {obs }}$, indicative of a flatter submm slope, have already been noted with Herschel by, e.g., O'Halloran et al. (2010); Grossi et al. (2010); Boselli et al. (2012) for subsolar-metallicity galaxies. This flatter slope may be the sign of a contribution from an extra emission in excess of the commonly used $\beta_{\text {theo }}=2$ models. Thus the flattening of the observed submm slope $\left(\beta_{o b s}<2\right)$ could be used as a diagnosis for possible excess emission appearing at $500 \mu \mathrm{~m}$ (see Section 5.4).

### 5.2 FIR/submm modelling of the cold dust properties

To complete our observational and qualitative view of the FIR-submm behaviour of the DGS and KINGFISH galaxies, we use a modified blackbody model to quantitatively determine the parameters discussed above: $T$ and $\beta_{o b s}$, as well as $M_{d u s t, B B}$ and $L_{F I R}$ in the DGS and KINGFISH samples. In the following, we note the dust mass estimated from a modified blackbody $M_{d u s t, B B}$, as opposed to the dust mass derived from a semi-empirical dust model, $M_{\text {dust }}$ (see Chapter 7).

### 5.2.1 Modelling the cold dust

Here we are considering Herschel data covering in the FIR-submm, where the bulk of the IR luminosity is emitted and where most of the mass of the cold dust can be traced. We know that cold grains are also large grains and thus the dust grains are likely to be in equilibrium with the
radiation field. If we assume that all of the grains have the same equilibrium temperature (i.e., the same size), their emission can be represented by a modified blackbody.

A single modified blackbody is thus fitted to the Herschel data of each galaxy from both DGS and KINGFISH samples and the modelled flux densities are given by:

$$
\begin{equation*}
F_{\nu}(\lambda)=\frac{M_{d u s t, B B} \kappa\left(\lambda_{0}\right)}{D^{2}}\left(\frac{\lambda}{\lambda_{0}}\right)^{-\beta_{o b s}} B_{\nu}(\lambda, T) \tag{5.1}
\end{equation*}
$$

where $\kappa\left(\lambda_{0}\right)=4.5 \mathrm{~m}^{2} \mathrm{~kg}^{-1}$ is the dust mass absorption opacity at the reference wavelength, $\lambda_{0}=100 \mu \mathrm{~m}$ (see Eq. 1.13 from Section 1.2.3). The free parameters are the temperature ( $T$ ) and dust mass ( $M_{\text {dust }, B B}$ ) as well as the emissivity index ( $\beta_{\text {obs }}$ ), where we leave $\beta_{o b s}$ free in the $[0.0,2.5]$ range. $D$ is the distance to the source (given in Table 3.5) and $B_{\nu}(\lambda, T)$ is the Planck function. Colour corrections, described in Sections 3.2.1 and 3.2.2, are included in the fitting procedure.
$\kappa\left(\lambda_{0}\right)$ has been calculated from the grain properties of Zubko et al. (2004), as in Galliano et al. $(2011)^{1}$, and is consistent with a $\beta_{\text {theo }}=2$ (see Section 7.1.2 and Fig. 7.3). Leaving $\beta_{\text {obs }}$ to vary in our fit can produce lower dust masses for lower $\beta_{o b s}$ (Bianchi 2013). This effect is discussed for the two dust masses relations we derive in Section 5.3. Moreover, this particular choice for the value $\kappa\left(\lambda_{0}\right)$ will only affect the absolute values of the dust masses. Choosing another model to derive $\kappa\left(\lambda_{0}\right)$ would not affect the intrinsic variations noted in Section 5.3.

At $70 \mu \mathrm{~m}$, possible contamination by dust grains that are not in thermal equilibrium, and whose emission cannot be represented by a modified blackbody, can occur in galaxies. An excess at $70 \mu \mathrm{~m}$ compared to a modified blackbody model can also appear, as seen in Fig. 5.2, because dust grains in a galaxy are more likely to have a temperature distribution rather than a single temperature. For example in spiral galaxies (present in the KINGFISH sample), the dust emission at $70-500 \mu \mathrm{~m}$ can originate from two components with different heating sources and potentially different temperatures (Bendo et al. 2010a, 2012; Boquien et al. 2011; Smith et al. 2012b). Therefore, we restrict our wavelength fitting range to $100-500 \mu \mathrm{~m}$. The $70 \mu \mathrm{~m}$ observation can be useful as an upper limit for a single temperature dust component. We redo the fit including the $70 \mu \mathrm{~m}$ point only if the modelled point from the fit without $70 \mu \mathrm{~m}$ data violates this upper limit condition, i.e., if it is greater than the observed point (e.g., Mrk 209 in Figure 5.4).

Some of our galaxies are not detected at some wavelengths. To have enough constraints for the fit, at least a detection up to $250 \mu \mathrm{~m}$ is required. If the galaxy is not detected beyond $160 \mu \mathrm{~m}$, we fit a modified blackbody including the $70 \mu \mathrm{~m}$ point. Indeed some galaxies peaking at very short wavelengths have their Rayleigh Jeans contribution already dropping at FIR and submm wavelengths, and are often not detected by SPIRE. For these galaxies, the $70 \mu \mathrm{~m}$ point is already on the Rayleigh Jeans side of the modified blackbody, and in this case we also include it in our fit.

All of these conditions are matched for 35 DGS galaxies, and we use the $70 \mu \mathrm{~m}$ point for 11 of them (five because of the violation of the upper limit condition at $70 \mu \mathrm{~m}$, four because the galaxy is not detected beyond $160 \mu \mathrm{~m}$, and two because the galaxy is not observed by SPIRE, see Table 5.1 for details). However, none of the 56 galaxies from KINGFISH requires using the $70 \mu \mathrm{~m}$ point.

From the fitted modified blackbodies, we also derive the FIR luminosity, $L_{F I R}$, by integrating the modelled curve between 50 and $650 \mu \mathrm{~m}$. The resulting parameters from the fits are given in Table 5.1. The SEDs are shown in Fig. 5.13 for all 35 DGS galaxies and Fig. 5.14 for the 56 KINGFISH galaxies.

### 5.2.2 Rigorous error estimation

In order to derive conservative errors for our $T, \beta, M_{d u s t, B B}$ and $L_{F I R}$ parameters we performed Monte Carlo iterations for each fit, following the method in Galliano et al. (2011). For each galaxy,

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Fig. 5.4. Examples of modified blackbody fits: the observed points are the red crosses whereas the modelled points are the filled blue circles. Upper limits are indicated with red diamonds. The bottom panel of each plot indicates the residuals from the fit. (top) Fit for Haro3, the observed $70 \mu \mathrm{~m}$ point which is not considered at first in our fitting procedure, is above the modelled one. (centre) Fit for Mrk209. Here the observed 70 $\mu \mathrm{m}$ point is below the modelled one, and the fit should be redone, giving us: (bottom) Fit for Mrk209 using the $70 \mu \mathrm{~m}$ point. Note how the shape of the modified blackbody varies between the two: for example, thy dust temperature for Mrk209 goes from 56 K (without $70 \mu \mathrm{~m}$ ) to 33 K (with $70 \mu \mathrm{~m}$ ).
we randomly perturb our fluxes within the error bars and perform fits of the perturbed SEDs (300 for each galaxy). To be able to do so, we must first carefully identify the various types of error and take special care for errors which are correlated between different bands.

As explained in Sections 4.2.2 and 4.3.2, we have measurement errors and calibration errors in our error estimates. The measurement errors are independent from one band to another and are usually well represented by a Gaussian distribution. The calibration errors, however, are correlated between different bands as it is the error on the flux conversion factor. It can be summarized for our case as follow:

PACS: Although the total calibration error is $5 \%$ in the three PACS bands it can be decomposed into two components:

- the uncertainty on the calibration model is $5 \%$ (according to the PACS photometer pointsource flux calibration documentation ${ }^{2}$ ) and is correlated between the three bands.
- the uncertainties due to noise in the calibration observations are: $1.4,1.6,3.5 \%$ at 70,100 , $160 \mu \mathrm{~m}$, respectively (PACS photometer point-source flux calibration). These uncertainties are independent from one band to another.
SPIRE: The SPIRE ICC recommend using $7 \%$ in each band but here again we can decompose it:
- the uncertainty on the calibration model is $5 \%$ (SPIRE Observer's Manual) and is correlated between the three bands.
- the uncertainties due to noise in the calibration observations are $2 \%$ for each band (SPIRE Observer's Manual). These uncertainties are independent.
- As SPIRE maps are given in Jy beam ${ }^{-1}$, the error on the beam area will also affect the calibration. The uncertainty on the beam area is given to be $4 \%$ in each band ${ }^{3}$ and are independent. This error is applied only to the extended sources, as the photometry for the point sources has been done directly on the timeline data.

The perturbation of the observed fluxes will then be the sum of two components:

- A normal random independent variable representing the measurement errors.
- A normal random variable describing the calibration errors that takes into account the correlation between the wavebands as described above, the same for each galaxy.

After performing 300 Monte-Carlo iterations, a distribution for each of the three model parameters $T, \beta, M_{d u s t, B B}$ as well as for $L_{F I R}$ is obtained for each galaxy (see example on Fig. 5.5). We chose to quote the $66.67 \%$ confidence level for our parameters defined by the range of the parameter values between 0.1667 and 0.8333 of the repartition function. As the distributions are often asymmetric we obtain asymmetric error bars on our parameters. These error bars are given in Table 5.1.

### 5.3 Dust FIR properties

We now have the $T, \beta, M_{d u s t, B B}$ and $L_{F I R}$ distributions of the DGS and KINGFISH samples (Figures 5.6 and 5.7). Note that KINGFISH is not a volume- or flux-limited sample but a crosssection of galaxies with different properties. Due to the heterogeneity of both samples we thus quote the median rather than the mean to compare the samples.

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Fig. 5.5. Examples of the obtained parameter distributions for the 300 Monte-Carlo iterations for the modified blackbody fits: (top) distribution of temperature, $T$, for Haro3, (bottom) distribution of dust mass, $M_{d u s t, B B}$, for Mrk209 with the $70 \mu \mathrm{~m}$ point included in the fit. The plain blue line notes the value of the parameter and the dashed blue lines note the $66.67 \%$ confidence level for the parameters.

### 5.3.1 Temperature

The range in dust temperature of the DGS galaxies is 21 to 98 K with a median $T \sim 32 \mathrm{~K}$ (Figure 5.6a). The most metal-poor galaxies are among the warmer ones. If we compare the KINGFISH to the DGS galaxies, our lowest temperatures are quite comparable ( 17 vs 21 K ), but the DGS galaxies have higher maximal dust temperatures ( 39 vs 98 K ). In Figure 5.6a, we see that the KINGFISH dust temperature distribution has a narrow peak around $\sim 20-25 \mathrm{~K}$ whereas the DGS distribution is broader. This difference is due to some galaxies in our sample that peak at extremely short wavelengths, a distinguishing feature of star-forming dwarf galaxies, resulting in very high dust temperatures for a single modified blackbody fit. The dust in DGS galaxies is thus overally warmer than that in more metal-rich galaxies $\left(T_{D G S}^{m e d}=32 \mathrm{~K}\right.$ and $\left.T_{K I N G F I S H}^{m e d}=23 \mathrm{~K}\right)$. This is coherent with the temperature trends presented in the previous section. Note that the high temperature tail of the DGS temperature distribution could be even more prominent than presented here: some galaxies are not detected beyond $100-160 \mu \mathrm{~m}$ rendering impossible the determination of their dust temperature with a modified blackbody fit (13 galaxies in the DGS). The SEDs for these galaxies likely peak at very short wavelengths giving very warm averaged dust temperatures (see Chapter 7 for the complete SED modelling of these galaxies).

### 5.3.2 Emissivity index

The "observed" emissivity index ( $\beta_{\text {obs }}$, see Eq. 5.1) distribution is shown on Fig. 5.6b, spanning a range from 0.0 to 2.5 with a median $\beta_{o b s} \sim 1.7$ for the DGS. There does not appear to be any clear correlation with metallicity here. Nonetheless, even if some DGS galaxies are nicely fitted by an often-presumed $\beta_{o b s}=2.0$ blackbody, some require a $\beta_{o b s} \leq 2.0$, and those are primarily metal-poor to moderately metal-poor galaxies ( 0.10 to $0.4 Z_{\odot}$ ). Note also that for the KINGFISH sample, all of the galaxies, but two, within this metallicity range have $0.5 \leq \beta_{o b s} \leq 2.0$. From SPIRE band ratios,

Boselli et al. (2012) also found that low-metallicity galaxies from the HRS sample were presenting submm colours consistent with an emissivity index $\leq 2.0$. Arbitrarily fixing $\beta=2.0$ in a blackbody modelling, in order to mimic the emissivity index appropriate for a mixture of amorphous silicate and graphite (reproducing the Milky Way observations), may not always be appropriate for lowmetallicity galaxies. However, we note that several DGS galaxies suggest a $\beta_{o b s}=0.0-0.5$ from our modified blackbody model (Fig. 5.6b). These six galaxies with $\beta_{o b s}<0.5$ in the DGS, are not detected beyond $160 \mu \mathrm{~m}$ and such a low $\beta_{\text {obs }}$, physically unrealistic, is probably due to the poorly constrained submm part of the SED.

In summary, there are metal-poor to moderately metal-poor galaxies, with metallicities between 0.1 and $0.4 Z_{\odot}$, for which $0.5 \leq \beta_{o b s} \leq 2.0$. These lower $\beta_{o b s}$ values, not necessarily realistic in term of actual grain properties, are representative of a flatter submm slope in the FIR observations, and could perhaps be an indicator of the presence of a submm excess in these sources (see Section 5.4).

### 5.3.3 Dust mass

The dust masses estimated from our modified blackbody fits for the DGS range from $1.0 \times 10^{2} M_{\odot}$ to $2.5 \times 10^{7} M_{\odot}$ (Figure 5.6 c ), with a median of $\sim 1.2 \times 10^{5} M_{\odot}$. From Figure 5.6 c we see that the most metal-poor galaxies are the least massive galaxies compared to the moderately metal-poor galaxies. The dwarf galaxies are, not surprisingly, less massive in dust than the galaxies from the KINGFISH sample: the median dust mass of the KINGFISH sample is about two orders of magnitude higher than for the DGS: $\sim 1.1 \times 10^{7} M_{\odot}$. In order to determine if this is only an effect due to the smaller sizes of dwarfs, we consider the ratio between the dust and stellar masses. The stellar masses for the DGS and KINGFISH samples have been computed in Chapter 3. Figure 5.7 shows that there is a strong decrease (about an order of magnitude) of the proportion of dust mass relative to the stellar mass with decreasing metallicity: we have a Spearman rank coefficient ${ }^{4} \rho=0.51$. The median for the ratio $M_{d u s t, B B} / M_{\star}$ is $0.02 \%$ for DGS versus $0.04 \%$ for KINGFISH. The best power-law fit gives:

$$
\begin{equation*}
M_{\text {dust }, B B} / M_{\star}=1.5 \times 10^{-20} \times(12+\log (O / H))^{18.4} \tag{5.2}
\end{equation*}
$$

However the dust masses derived here for both samples are probably lower limits of the real dust masses in many cases (see Dale et al. 2012 and Section 8.2.1). Indeed we allow our $\beta_{\text {obs }}$ to go to very low values, giving lower dust masses than if we fixed it to 1.5 or even 2.0 : as we allow a greater emission effciency for the grains, we need less mass than if we were using a higher emissivity index, to account for the same amount of luminosity. We perform the test by fixing the emissivity index parameter to 2.0 and find that the correlation between the dust-to-stellar mass ratios with metallicity almost vanishes: with $\beta=2.0$ modified blackbody dust masses, the Spearman rank coefficient for the relation decreases to $\rho=0.08$. Additionally, with our modified blackbody fits we are considering only one temperature and grain size. We may be missing here a fraction of the dust mass coming from warmer big grains, and this contribution may be more important in low-metallicity galaxies rather than in more metal-rich ones. The mass corresponding to the stochastically heated grains is, however, negligible. In Chapter 7, we will obtain total dust masses from a full semi-empirical SED model, which will allow us to study the dust-to-stellar mass ratios in more details.

### 5.3.4 FIR luminosity

The FIR luminosities in the DGS sample range from $1.2 \times 10^{7} L_{\odot}$ to $5.3 \times 10^{10} L_{\odot}$ (Figure 5.6d), with a median of $\sim 5.3 \times 10^{8} L_{\odot}$. We see in Figure 5.6 d that dwarf galaxies are less luminous in the

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Fig. 5.6. Distributions of temperature (a), emissivity index (b), dust mass (c) and FIR luminosity (d) from modified blackbody fits for Herschel data for the DGS and KINGFISH samples. The colour scale represents the range of metallicity values. On each panel, the upper/lower histogram is the KINGFISH/DGS distribution for the parameter.


Fig. 5.7. (top) $M_{d u s t, B B} / M_{\star}$ as a function of metallicity for DGS (purple crosses) and KINGFISH (orange downward triangles). The best power-law fit is indicated as a black line, and corresponds to: $\log \left(M_{d u s t, B B} / M_{\star}\right)=(-20.4 \pm 1.6)+(18.4 \pm 1.8) \times \log (12+\log (\mathrm{O} / \mathrm{H}))$. The distribution of $M_{d u s t, B B} / M_{\star}$ is indicated on the side for both samples: plain purple line for DGS and dashed orange line for KINGFISH. (bottom) $L_{F I R} / M_{d u s t, B B}$ as a function of metallicity for DGS (crosses) and KINGFISH (downward triangles). The colours code the temperature, $T$. The best power-law fit line is indicated as a black line, and corresponds to: $\log \left(L_{F I R} / M_{d u s t, B B}\right)=(24.4 \pm 1.1)+(-23.6 \pm 1.2) \times \log (12+\log (\mathrm{O} / \mathrm{H}))$. The distribution of $L_{F I R} / M_{d u s t, B B}$ is indicated on the side for both samples: solid line for DGS and dashed line for KINGFISH. On both plots: the errors on the metallicities are omitted for clarity. They are of about 0.1 dex on average.

FIR than the galaxies from the KINGFISH sample. However if we consider $L_{F I R} / M_{d u s t, B B}$, which represents the quantity of light emitted by the available amount of dust, there is a strong trend of increasing $L_{F I R} / M_{d u s t, B B}$ with decreasing metallicity (Figure 5.7): here we have a Spearman rank coefficient $\rho=-0.72$. The best power-law fit gives:

$$
\begin{equation*}
L_{F I R} / M_{\text {dust }, B B}=4.2 \times 10^{24} \times(12+\log (O / H))^{-23.6} \tag{5.3}
\end{equation*}
$$

Despite their lower dust masses, dwarf galaxies emit more in the FIR/submm than more metalrich galaxies, per unit dust mass ( $\sim 6$ times more for the DGS). Here fixing $\beta_{o b s}=2$ only decreases the strength of the correlation $(\rho=-0.46)$ but the correlation still holds. This difference in $L_{F I R} / M_{d u s t, B B}$ between the two samples is a direct consequence of the higher temperature of dust grains in dwarf galaxies, as shown by the colours on Fig. 5.7, due to the stronger and harder ISRF in which the grains are embedded. However, as mentioned above, the total dust mass may be underestimated by the modified blackbody model in lower metallicity galaxies and this trend could be weaker. This will be explored in Chapter 8.

### 5.3.5 Temperature - emissivity index relation



Fig. 5.8. Temperature versus $\beta_{o b s}$ from the modified blackbody fits for the DGS (purple crosses) and for KINGFISH galaxies (orange downward triangles). The dashed lines correspond to the best power-law fit for the DGS (purple line) and KINGFISH (orange line) galaxies, excluding the galaxies for which $\beta_{o b s}=0.0$. They correspond to: $\log \left(T_{D G S}\right)=(1.59 \pm 0.01)+(-0.48 \pm 0.04) \times \log \left(\beta_{o b s, D G S}\right)$, and $\log \left(T_{\text {KINGFISH }}\right)=$ $(1.41 \pm 0.02)+(-0.29 \pm 0.05) \times \log \left(\beta_{o b s, \text { KINGFISH }}\right)$. For comparison, the analysis of Dupac et al. (2003) has been overplotted on our data (black dotted line). For clarity the error bars on the parameters have been displayed only for the DGS sample.

Some studies have noted an inverse $\beta$ / temperature correlation in objects from starless cores to galaxies (Dupac et al. 2003; Yang \& Phillips 2007; Anderson et al. 2010; Paradis et al. 2010; Planck Collaboration et al. 2011a; Galametz et al. 2012; Smith et al. 2012a). To investigate this possible effect in our samples, we plot these two parameters from our modified blackbody fits ( $T$ and $\beta_{o b s}$ ) (Figure 5.8).

First we note that the DGS galaxies have overall higher dust temperature than the KINGFISH galaxies as already noted in Figures 5.2, 5.3 and 5.6. We also have the DGS galaxies where the fit gives $\beta_{o b s}=0.0$, without detections beyond $160 \mu \mathrm{~m}$ that we believe to be due to a poorly constrained submm SED. If we exclude these galaxies, the KINGFISH and DGS samples present an anticorrelation between $T$ and $\beta_{o b s}$, and this anticorrelation seems to be steeper in the DGS: the best power-law fit gives $T \propto \beta_{o b s}^{-0.48}$ for the DGS and $T \propto \beta_{o b s}^{-0.29}$ for the KINGFISH galaxies. However, the anticorrelation seems stronger in KINGFISH than in the DGS sample $\left(\rho_{\text {KINGFISH }}=-0.69 \mathrm{vs}\right.$ $\left.\rho_{D G S}=-0.56\right)$.

Shetty et al. (2009a,b) and Juvela \& Ysard (2012a,b) showed that such an observed anticorrelation comes from the assumption of a constant temperature along the line-of-sight in modified blackbody fits and from noise in the measurements. They advise caution when interpreting this $\beta /$ temperature relationship when derived from $\chi^{2}$ modified blackbody fits. Kelly et al. (2012) show that a $\chi^{2}$ fit can artificially produce an anticorrelation between $T$ and $\beta_{o b s}$, whereas a Bayesian fit does not, and recovers the true parameters more accurately.

Nonetheless, if we assume that the differences in the observed ( $T, \beta_{o b s}$ ) relations between DGS and KINGFISH can be due to changes in dust optical grain properties in the submm (as suggested by Meny et al. 2007; Paradis et al. 2010), this may be the sign that the assumption of a single grain temperature, the presence of noise in the measurements and the use of a $\chi^{2}$ fitting procedure may only be partially responsible for the observed trends. However, given the very large errors on the $T$ and $\beta$ parameters, it is difficult to draw a solid conclusion on this issue.

### 5.4 The intriguing submm excess: the modified blackbody view

A submm excess has been observed in the past in several low-metallicity galaxies (Galliano et al. 2003, 2005; Dumke et al. 2004; Bendo et al. 2006; Zhu et al. 2009; Galametz et al. 2009; Bot et al. 2010; Grossi et al. 2010; Galametz et al. 2011; Dale et al. 2012). It has been called "excess" because the current available models are unable to fully explain the submm emission of these galaxies. This submm excess has been one of the main sources of uncertainty in dust modelling in dwarf galaxies for the past few years, especially on the dust mass parameter.

### 5.4.1 Characterisation of the excess

In most models, $\beta_{\text {theo }}=2$ is often assumed in order to mimic the optical properties of the dust grain mixture of the Galaxy. In spiral galaxies, a modified blackbody with a fixed $\beta_{\text {theo }}$ to 2 reproduces well the FIR emission (Bendo et al. 2003, 2010a). In the colour-colour diagrams we hinted that a low $\beta_{\text {obs }}$ may be the sign of a possible presence of an excess emission adding its contribution to a $\beta_{\text {theo }}=2$ submm SED. Boselli et al. (2012) also showed that the $F_{250} / F_{500}$ colour was more consistent with an effective emissivity index of 1.5 for the lowest metallicity galaxies in the HRS sample. Here we want to determine, systematically, which galaxies of the DGS and KINGFISH samples present an excess. A modified blackbody with a fixed emissivity index $\beta_{\text {theo }}$ of 2.0 is fit to the data for both DGS and KINGFISH samples. Here again, we use the $70 \mu \mathrm{~m}$ point only if the modelled flux is larger than the observed flux.

We take the relative residual at $500 \mu \mathrm{~m}$ to be:

$$
\begin{equation*}
R(500)=\frac{L_{\nu}^{\text {observed }}(500)-L_{\nu}^{\text {model }}(500)}{L_{\nu}^{\text {model }}(500)} \tag{5.4}
\end{equation*}
$$

In order to define a residual, $R$, at $500 \mu \mathrm{~m}$ the galaxy must be detected out to $500 \mu \mathrm{~m}$. This, unfortunately, reduces our sample to 78 galaxies due to the high number of faint galaxies in the DGS sample.

Following the same procedure as in Section 5.2.2, we randomly perturb the fluxes within the errors bars and perform fits of the perturbed SEDs (300 for each galaxy). A distribution of $R(500)$ is generated and the $66.67 \%$ confidence level of the distribution gives the error on the residual at $500 \mu \mathrm{~m}: \Delta R(500)$. The values of $R(500)$ and $\Delta R(500)$ are listed in Table 5.3.

A galaxy is then flagged with "excess" if the relative residual at $500 \mu \mathrm{~m}$ is greater than the corresponding error: $R(500)>\Delta R(500)$ (see Table 5.3). As the $500 \mu \mathrm{~m}$ point is included in the fit, the procedure will also try to achieve a good fit of the $500 \mu \mathrm{~m}$ point, and this will give lower $R(500)$ than if the $500 \mu \mathrm{~m}$ point was not included in the fit. That is why we fix our "excess" criterion to a $1 \sigma$ detection only. For both samples, the $R(500)$ distribution is shown in Figure 5.9, and excess galaxies are indicated by hashed cells.


Fig. 5.9. Relative residual distribution at $500 \mu \mathrm{~m}$ for modified blackbody fits with a fixed $\beta_{\text {theo }}$ of 2.0 for DGS (purple) and KINGFISH (orange) samples. Galaxies for which the residual at $500 \mu \mathrm{~m}$ is greater than the corresponding error bar $(R(500)>\Delta R(500))$ have been marked by hashed cells.

Out of 78 galaxies, $42 \%$ presents an excess at $500 \mu \mathrm{~m}$ with respect to a $\beta_{\text {theo }}=2$ modified blackbody: nine are from DGS and 24 from KINGFISH. It is interesting to note that seven out of the nine KINGFISH galaxies of Irregular type ( $\mathrm{Im}, \mathrm{I} 0$ or Sm ) detected at $500 \mu \mathrm{~m}$, are among the 24 "excess" KINGFISH galaxies. The ones missing are M81dwB and HoII which both have a very large error bar on the $500 \mu \mathrm{~m}$ flux and thus a very wide $R(500)$ distribution. Dale et al. (2012) looked at the residual at $500 \mu \mathrm{~m}$ for a Draine \& Li (2007) model fit (see Dale et al. 2012 for details) and also found that most of these Irregular galaxies presented an excess at $500 \mu \mathrm{~m}$. They mention a dozen KINGFISH galaxies with a $R(500)$ above $60 \%$. However as their study is based on a different model than ours, we will not go deeper into any further comparison.

Figure 5.10 shows the metallicity distribution of the 33 excess galaxies (black line) together with the joint metallicity distribution of DGS and KINGFISH samples (grey line). Note how the absence of detections at $500 \mu \mathrm{~m}$ reduces the low-metallicity tail of the joint metallicity distribution.

The metallicity distribution for the excess galaxies peaks around $12+\log (\mathrm{O} / \mathrm{H}) \sim 8.3$ and is skewed towards the low-metallicity end of the distribution: $64 \%$ of the excess galaxies are galaxies with $Z<0.4 Z_{\odot}$, whereas, in the total distribution of galaxies detected at $500 \mu \mathrm{~m}$, (grey line on Fig. 5.10), only $46 \%$ of the total number of galaxies are galaxies with $Z<0.4 Z_{\odot}$. Moreover, the proportion of excess galaxies in the [7.5-8.3] metallicity range is $\sim 58 \%$ versus $\sim 29 \%$ in the $] 8.3-8.8$ ] range. This shows that the submm excess seems to occur mainly in metal-poor galaxies, at least when a $\beta_{\text {theo }}=2$ modified blackbody model is used. The colours on Fig. 5.10 code the signal-to-noise ratio of the residual at $500 \mu \mathrm{~m}$ for the excess galaxies. There seems to be a dichotomy in the distribution around $12+\log (\mathrm{O} / \mathrm{H}) \sim 8.3$, with the strongest excesses being detected in the lowest metallicity galaxies.


Fig. 5.10. Metallicity distribution for the excess galaxies in black. The colours mark the presence of an excess at $500 \mu \mathrm{~m}$ and code the intensity of this excess: $R(500) / \Delta R(500)$. The metallicity distribution for the DGS \& KINGFISH galaxies detected at $500 \mu \mathrm{~m}$ is outlined in grey. The grey cells mark all of the galaxies for which no detection is available at $500 \mu \mathrm{~m}$.

On Fig. 5.11, we have a clear anti-correlation between $R(500)$ and $\beta_{\text {obs }}$ from the modified blackbody fits from Section 5.3: $\rho=-0.78$. For galaxies with a "naturally" flatter slope (i.e., a low $\beta_{o b s}$ ), forcing a steeper slope (i.e., fixing $\beta_{\text {theo }}=2.0$ ) will naturally increase the residuals at the longest wavelengths, thus generating the correlation between $R(500)$ and $\beta_{o b s}$.

All of the galaxies showing an excess (i.e., $R(500)>\Delta R(500)$ ) have indeed a low $\beta_{o b s}\left(\beta_{o b s} \leq 2.0\right)$ (Fig. 5.11). It is also interesting to note that this corresponds to $80 \%$ of the 44 galaxies with $\beta_{o b s} \leq$ 2.0. On the colour-colour diagram of Fig. 5.3, all of the excess galaxies fall on the left side of the $\beta_{\text {theo }}=2.0$ line. Moreover, all of the galaxies falling on the left side of the $\beta_{\text {theo }}=1.5$ line, except one, present an excess. This is coherent with what is observed on Fig. 5.11, and can be useful to select potential targets for FIR/submm follow-up observations.

### 5.4.2 A word of caution: submm excess appearing beyond $500 \mu \mathrm{~m}$

In the previous analysis we are considering only Herschel wavelengths, and any galaxy for which a submm excess is appearing beyond Herschel wavelengths would not be detected here. This is illustrated with two galaxies of the DGS sample with observations beyond $500 \mu \mathrm{~m}$, Haro 11 and II $Z_{w} 40$, both modelled with modified blackbodies, with the same procedure as in Section 5.4.1. Haro


Fig. 5.11. Relative residual at $500 \mu \mathrm{~m}$ versus $\beta_{o b s}$ (from Section 5.3) for DGS (crosses) and KINGFISH (downward triangles) galaxies. The green symbols mark the galaxies presenting an excess at $500 \mu \mathrm{~m}$. The best fit line is indicated as a solid black line, and corresponds to: $\mathrm{R}(500)=(66.6 \pm 7.1)+(-29.2 \pm 3.4) \times$ $\beta_{o b s}$.

11 falls to the left side of the $\beta=1.5$ line and has been identified as an "excess" galaxy in Table 5.3 whereas II Zw 40 falls to the right side of the $\beta=1.5$ line (Figure 5.3) and does not present any excess at $500 \mu \mathrm{~m}$ when using only Herschel bands (Table 5.3).

As shown in Figure 5.12, Haro 11 presents an excess at $500 \mu \mathrm{~m}(R(500) \sim 28 \% \pm 17 \%)$, confirmed at $870 \mu \mathrm{~m}(R(870) \sim 360 \% \pm 15 \%)$. However, the submm excess is clearly appearing at longer wavelengths ( $\geq 500 \mu \mathrm{~m}$ ) for II Zw 40 when including observations beyond $500 \mu \mathrm{~m}$. At $500 \mu \mathrm{~m}$ $R(500) \sim 2 \pm 8 \%$, and $R(450) \sim 6 \pm 34 \%$ but at 850 and $1200 \mu \mathrm{~m}$ we have $R(850) \sim 265 \pm 16 \%$ and $R(1200) \sim 370 \pm 40 \%$. This illustrates the need for submm data, to complement the existing Herschel data.

### 5.5 Limitations of the modified blackbody approach

Throughout this Chapter we saw that the modified blackbody modeling presents some limitations due to the very strong assumption of a single temperature for the dust grains. Although this approach, popular in the literature, provided useful constraints on the dust properties, a more realistic dust model is now needed to really pinpoint the dust properties, and their variations with metallicity (i.e., dust-to-stellar mass and (F)IR luminosity-to-dust mass ratios). A semi-empirical dust SED model is thus used in Part III and now allows the grains to vary in sizes (and thus temperatures). The necessary ancillary data is collected for the two samples and presented in Chapter 6. The dust model is presented in Chapter 7 and the dust properties analysed in Chapter 8.

Table 5.1. Table of modified blackbody fit parameters for the DGS galaxies.

| Source | Temperature (K) | $\beta_{\text {obs }}$ | $\mathrm{M}_{\text {dust }, B B}\left(M_{\odot}\right)$ | $\mathrm{L}_{F I R-B B}\left(L_{\odot}\right)$ |
| :---: | :---: | :---: | :---: | :---: |
| Haro11 | $38_{-6}^{+11}$ | $1.96{ }_{-0.41}^{+0.44}$ | $5.0_{-3.3}^{+6.5} \times 10^{6}$ | $5.3{ }_{-0.5}^{+1.1} \times 10^{10}$ |
| Haro2 | $25_{-1}^{+3}$ | $2.388_{-0.38}^{+0.09}$ | $2.1_{-0.9}^{+3.5} \times 10^{6}$ | $2.5_{-0.1}^{+0.1} \times 10^{9}$ |
| Haro3 | $26_{-2}^{+3}$ | $2.15{ }_{-0.34}^{+0.31}$ | $1.7_{-0.8}^{+1.1} \times 10^{6}$ | $2.4_{-0.1}^{+0.2} \times 10^{9}$ |
| He2-10 | $26_{-1}^{+4}$ | $2.24_{-0.40}^{+0.21}$ | $1.3{ }_{-0.7}^{+0.5} \times 10^{6}$ | $2.0_{-0.1}^{+0.1} \times 10^{9}$ |
| HS0017+1055 ${ }^{2 a}$ | $98_{-36}^{+34}$ | $0.00_{-0.00}^{+1.34}$ | $1.9_{-1.4}^{+8.2} \times 10^{3}$ | $3.4_{-0.4}^{+0.3} \times 10^{8}$ |
| HS0052+2536 | $37_{-9}^{+14}$ | $1.20_{-0.73}^{+0.79}$ | $1.1_{-0.8}^{+3.1} \times 10^{6}$ | $8.5_{-1.0}^{+1.6} \times 10^{9}$ |
| HS0822+3542 | - | - | - | - |
| HS1222+3741 | - | - | - | - |
| HS1236+3937 | - | - | - | - |
| HS1304+3529 ${ }^{1}$ | $32_{-3}^{+7}$ | $2.05_{-0.72}^{+0.47}$ | $2.2{ }_{-1.3}^{+1.5} \times 10^{5}$ | $9.4_{-0.4}^{+0.5} \times 10^{8}$ |
| HS1319+3224 | - | - | ${ }_{-}$ | ${ }_{-}$ |
| HS1330 $+3651^{2 b}$ | $50_{-8}^{+3}$ | $0.00_{-0.00}^{+0.61}$ | $3.3{ }_{-0.6}^{+3.4} \times 10^{4}$ | $8.8{ }_{-0.4}^{+0.4} \times 10^{8}$ |
| HS1442+4250 | - | - | - | - |
| HS2352+2733 | - | - | - | - |
| IZw18 | - | - | - | - |
| IC10 | $21_{-1}^{+3}$ | $2.25_{-0.49}^{+0.26}$ | $2.6_{-1.4}^{+1.1} \times 10^{5}$ | $1.1_{-0.05}^{+0.05} \times 10^{8}$ |
| IIZw40 | $33_{-4}^{+5}$ | $1.711_{-0.32}^{+0.41}$ | $1.9{ }_{-0.9}^{+1.8} \times 10^{5}$ | $9.2_{-0.7}^{+0.9} \times 10^{8}$ |
| Mrk1089 | $23_{-1}^{+3}$ | $2.34_{-0.39}^{+0.15}$ | $2.5{ }_{-1.2}^{+0.7} \times 10^{7}$ | $1.6_{-0.1}^{+0.1} \times 10^{10}$ |
| Mrk1450 | $43_{-13}^{+40}$ | $1.35{ }_{-0.92}^{+1.00}$ | $7.6_{-6.1}^{+24.6} \times 10^{3}$ | $1.22_{-0.2}^{+0.4} \times 10^{8}$ |
| Mrk153 ${ }^{1}$ | $32_{-2}^{+5}$ | $2.33_{-0.61}^{+0.15}$ | $1.2{ }_{-0.6}^{+0.4} \times 10^{5}$ | $5.3_{-0.2}^{+0.3} \times 10^{8}$ |
| Mrk209 ${ }^{1}$ | $34_{-3}^{+6}$ | $1.95_{-0.47}^{+0.42}$ | $2.1_{-1.1}^{+1.4} \times 10^{3}$ | $1.3{ }_{-0.1}^{+0.1} \times 10^{7}$ |
| Mrk930 | $26_{-2}^{+4}$ | $2.22_{-0.43}^{+0.21}$ | $5.7_{-3.0}^{+3.0} \times 10^{6}$ | $8.6_{-0.5}^{+0.6} \times 10^{9}$ |
| NGC1140 | $23_{-1}^{+2}$ | $2.17_{-0.36}^{+0.31}$ | $3.0_{-1.5}^{+1.7} \times 10^{6}$ | $1.9_{-0.1}^{+0.1} \times 10^{9}$ |
| NGC1569 | $28_{-2}^{+4}$ | $2.20_{-0.38}^{+0.28}$ | $2.8{ }_{-1.4}^{+1.5} \times 10^{5}$ | $5.7_{-0.4}^{+0.5} \times 10^{8}$ |
| NGC1705 | $33_{-4}^{+5}$ | $1.16_{-0.28}^{+0.33}$ | $8.4_{-3.7}^{+1.7} \times 10^{3}$ | $4.1_{-0.3}^{+0.3} \times 10^{7}$ |
| NGC2366 | $39_{-4}^{+4}$ | $0.96_{-0.23}^{+0.22}$ | $6.8{ }_{-2.1}^{+2.9} \times 10^{3}$ | $7.3_{-0.2}^{+0.3} \times 10^{7}$ |
| NGC4214 | $26_{-3}^{+3}$ | $1.39_{-0.37}^{+0.37}$ | $2.0_{-1.0}^{+2.0} \times 10^{5}$ | $2.9{ }_{-0.1}^{+0.2} \times 10^{8}$ |
| NGC4449 | $22_{-1}^{+3}$ | $2.18{ }_{-0.41}^{+0.27}$ | $2.3{ }_{-1.2}^{+1.2} \times 10^{6}$ | $1.4_{-0.1}^{+0.1} \times 10^{9}$ |
| NGC4861 | $28_{-3}^{+3}$ | $1.36_{-0.32}^{+0.34}$ | $6.1_{-2.7}^{+5.2} \times 10^{4}$ | $1.3{ }_{-0.1}^{+0.1} \times 10^{8}$ |
| NGC5253 | $30_{-3}^{+5}$ | $1.86{ }_{-0.37}^{+0.40}$ | $1.9_{-1.0}^{+1.8} \times 10^{5}$ | $5.5{ }_{-0.4}^{+0.5} \times 10^{8}$ |
| NGC625 | $29_{-3}^{+4}$ | $1.33_{-0.30}^{+0.35}$ | $5.9_{-2.7}^{+5.9} \times 10^{4}$ | $1.5{ }_{-0.1}^{+0.1} \times 10^{8}$ |
| NGC6822 | $26_{-4}^{+4}$ | $0.99_{-0.50}^{+0.59}$ | $1.1_{-0.6}^{+1.6} \times 10^{4}$ | $1.8{ }_{-0.1}^{+0.1} \times 10^{7}$ |
| Pox186 | $40_{-4}^{+4}$ | $0.00_{-0.00}^{+0.00}$ | $1.6_{-0.5}^{+0.7} \times 10^{3}$ | $2.0_{-0.2}^{+0.2} \times 10^{7}$ |
| SBS0335-052 ${ }^{2 a, 3}$ | $89_{-8}^{+10}$ | $1.64{ }_{-0.33}^{+0.39}$ | $8.0_{-1.4}^{+1.8} \times 10^{2}$ | $1.2{ }_{-1.2}^{+1.4} \times 10^{7}$ |
| SBS1159+545 | - | - | - | $-$ |
| SBS1211+540 ${ }^{2 a, 3}$ | $71_{-7}^{+7}$ | $0.34_{-0.41}^{+0.46}$ | $1.0_{-0.3}^{+0.4} \times 10^{2}$ | $1.2{ }_{-0.1}^{+0.1} \times 10^{7}$ |
| SBS1249+493 | - | - | $-$ | $-$ |
| SBS1415+437 ${ }^{2 b}$ | $35_{-3}^{+16}$ | $2.37_{-1.20}^{+0.15}$ | $4.9_{-3.1}^{+2.1} \times 10^{3}$ | $3.5_{-0.3}^{+0.3} \times 10^{7}$ |
| SBS1533+574 ${ }^{2 a}$ | $42_{-10}^{+8}$ | $0.44_{-0.40}^{+1.20}$ | $5.6_{-2.5}^{+11.4} \times 10^{4}$ | $8.2_{-0.4}^{+0.4} \times 10^{8}$ |
| Tol0618-402 | - | - | - | - |
| Tol1214-277 | - | - | - | - |
| UGC4483 | - | - | - | - |
| UGCA20 | - | - | - | - |
| UM133 | $41_{-15}^{+16}$ | $0.44_{-0.57}^{+1.89}$ | $3.1_{-2.8}^{+56.4} \times 10^{3}$ | $4.0_{-0.5}^{+0.9} \times 10^{7}$ |
| UM311 | $24_{-2}^{+3}$ | $1.58{ }_{-0.39}^{+0.34}$ | $3.7_{-1.9}^{+3.5} \times 10^{6}$ | $3.4_{-0.1}^{+0.2} \times 10^{9}$ |
| UM448 ${ }^{1}$ | $33_{-2}^{+2}$ | $1.99_{-0.16}^{+0.18}$ | $9.9{ }_{-2.5}^{+3.5} \times 10^{6}$ | $4.9_{-0.2}^{+0.2} \times 10^{10}$ |
| UM461 | $24_{-1}^{+1}$ | $2.50_{-0.00}^{+0.00}$ | $2.7_{-0.5}^{+0.5} \times 10^{4}$ | $2.5_{-0.2}^{+0.2} \times 10^{7}$ |
| VIIZw403 ${ }^{1}$ | $34_{-3}^{+3}$ | $1.57_{-0.23}^{+0.27}$ | $2.1_{-0.7}^{+1.1} \times 10^{3}$ | $1.2{ }_{-0.05}^{+0.1} \times 10^{7}$ |

1: $70 \mu \mathrm{~m}$ point included in fit: violation of the upper limit condition at $70 \mu \mathrm{~m}$.
${ }^{2 a}: 70 \mu \mathrm{~m}$ point included in fit: no detections beyond $160 \mu \mathrm{~m}$.
${ }^{2 b}$ : $70 \mu \mathrm{~m}$ point included in fit: no observations beyond $160 \mu \mathrm{~m}$.
${ }^{3}$ : For these particular galaxies, we included the $24 \mu \mathrm{~m}$ point in the fit as the $24 \mu \mathrm{~m}$ point fell below the modelled modifipe blackbody when we just overlaid it on the plot.

Table 5.2. Table of modified blackbody fit parameters for the KINGFISH galaxies.

| Source | Temperature (K) | $\beta_{\text {obs }}$ | $\mathrm{M}_{\text {dust, } B B}\left(M_{\odot}\right)$ | $\mathrm{L}_{F I R-B B}\left(L_{\odot}\right)$ |
| :---: | :---: | :---: | :---: | :---: |
| NGC0337 | $25 .{ }_{-3}^{+3}$. | $1.70_{-0.39}^{+0.46}$ | $6.0_{-2.6}^{+7.0} \times 10^{6}$ | $7.7_{-0.4}^{+0.4} \times 10^{9}$ |
| NGC0584 | - | - |  |  |
| NGC0628 | 19. ${ }_{-2}^{+2 .}$ | $2.19{ }_{-0.31}^{+0.54}$ | $2.0_{-0.5}^{+1.9} \times 10^{7}$ | $4.5{ }_{-0.2}^{+0.2} \times 10^{9}$ |
| NGC0855 | 25. ${ }_{-3}^{+3 .}$ | $1.411_{-0.37}^{+0.31}$ | $1.7_{-0.9}^{+1.5} \times 10^{5}$ | $2.1_{-0.1}^{+0.1} \times 10^{8}$ |
| NGC0925 | 23. ${ }_{-3}^{+3}$. | $1.10_{-0.41}^{+0.45}$ | $2.6{ }_{-1.2}^{+2.7} \times 10^{6}$ | $2.5_{-0.1}^{+0.1} \times 10^{9}$ |
| NGC1097 | $22 .{ }_{-2}^{+3}$ | $1.988_{-0.36}^{+0.46}$ | $4.3_{-1.8}^{+4.0} \times 10^{7}$ | $2.5{ }^{+0.1}+10^{+0.1}$ |
| NGC1266 | $26 .{ }_{-2}^{+4 .}$ | $2.31_{-0.42}^{+0.18}$ | $1.1_{-0.5}^{+1.4} \times 10^{7}$ | $1.5_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC1291 | $20 .{ }_{-2}^{+2}$. | $1.70_{-0.41}$ | $5.2_{-2.2}^{+4.4} \times 10^{6}$ | $1.7_{-0.1}^{+0.1} \times 10^{9}$ |
| NGC1316 | 20. ${ }_{-0.2}^{+1 .}$ | $2.50_{-0.18}^{+0.00}$ | $1.4{ }_{-0.2}^{+0.1} \times 10^{7}$ | $4.4{ }_{-0.2}^{+0.2} \times 10^{9}$ |
| NGC1377 | 34. | $1.79_{-0.43}^{+0.35}$ | $7.0_{-3.5}^{+5.1} \times 10^{5}$ | $3.9{ }_{-0.4}^{+0.5} \times 10^{9}$ |
| NGC1404 | - | - | $-$ |  |
| IC0342 | $20 .{ }_{-1}^{+3}$. | $2.17_{-0.46}^{+0.34}$ | $3.0_{-1.7}^{+1.6} \times 10^{7}$ | $1.0_{-0.04}^{+0.05} \times 10^{10}$ |
| NGC1482 | 24. ${ }_{-1}^{+4 .}$ | $2.36_{-0.49}^{+0.14}$ | $3.0_{-1.6}^{+0.9} \times 10^{7}$ | $2.6_{-0.1}^{+0.2} \times 10^{10}$ |
| NGC1512 | 25. ${ }_{-3}^{+3 .}$ | $0.988_{-0.35}^{+0.46}$ | $1.8{ }_{-0.8}^{+1.8} \times 10^{6}$ | $2.2_{-0.1}^{+0.1} \times 10^{9}$ |
| NGC2146 | $24 .{ }_{-1}^{+2}$. | $2.50_{-0.37}^{+0.00}$ | $7.3_{-2.3}^{+1.2} \times 10^{7}$ | $6.9_{-0.4}^{+0.5} \times 10^{10}$ |
| HoII | 25. ${ }_{-3}^{+9}$. | $1.888_{-0.89}^{-+0.35}$ | $3.2{ }_{-2.9}^{+5.6} \times 10^{4}$ | $3.6_{-0.4}^{+0.4} \times 10^{7}$ |
| DDO053 | - | - | $-$ | $-$ |
| NGC2798 | $26 .{ }_{-2}^{+4 .}$ | $2.24{ }_{-0.40}^{+0.28}$ | $1.3{ }_{-0.6}^{+0.8} \times 10^{7}$ | $1.8{ }_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC2841 | 17. ${ }_{-1}^{+2}$ | $2.33_{-0.57}^{+0.19}$ | $6.6{ }_{-37}^{+2.2} \times 10^{7}$ | $6.7_{-0.3}^{+0.3} \times 10^{9}$ |
| NGC2915 | 34. ${ }_{-6}^{10}$. | $0.73_{-0.41}^{+0.41}$ | $4.5{ }_{-2.0}^{+4.0} \times 10^{3}$ | $2.7_{-0.2}^{+0.3} \times 10^{7}$ |
| HoI | 39. ${ }_{-8}^{+7 .}$ | $0.00_{-0.00}^{\text {+0.48 }}$ | $7.0_{-3.1}^{+18.5} \times 10^{2}$ | $7.6_{-1.2}^{+0.4} \times 10^{6}$ |
| NGC2976 | 21. ${ }_{-2}$. | $1.99_{-0.46}^{+0.31}$ | $1.1_{-0.6}^{+0.7} \times 10^{6}$ | $4.9_{-0.2}^{+0.2} \times 10^{8}$ |
| NGC3049 | $32 .{ }_{-4}^{+6}$. | $0.82_{-0.36}^{+0.39}$ | $4.3{ }_{-2.4}^{+3.9} \times 10^{5}$ | $1.9_{-0.1}^{+0.1} \times 10^{9}$ |
| NGC3077 | $26 .{ }_{-3}^{+4 .}$ | $1.62_{-0.41}^{+0.43}$ | $3.1_{-1.5}^{+3.1} \times 10^{5}$ | $4.3{ }_{-0.2}^{+0.2} \times 10^{8}$ |
| M81dwB | $27 .{ }_{-7}^{+10}$. | $0.82_{-0.79}^{+1.09}$ | $1.4{ }_{-0.9}^{+6.9} \times 10^{3}$ | $3.0_{-0.4}^{+0.4} \times 10^{6}$ |
| NGC3190 | 18. ${ }_{-1}^{+2 .}$ | $2.50_{-0.52}^{+0.01}$ | $2.7_{-1.1}^{+0.4} \times 10^{7}$ | $4.5_{-0.2}^{+0.2} \times 10^{9}$ |
| NGC3184 | 19. ${ }_{-3}^{+1}$. | $2.18{ }_{-0.36}^{+0.51}$ | $2.7_{-0.7}^{+2.6} \times 10^{7}$ | $5.6_{-0.2}^{+0.2} \times 10^{9}$ |
| NGC3198 | 21. ${ }_{-2}^{+2}$. | $1.58{ }_{-0.39}^{+0.36}$ | $9.1_{-4.3}^{+10.2} \times 10^{6}$ | $4.8{ }_{-0.2}^{+0.2} \times 10^{9}$ |
| IC2574 | 27. ${ }_{-3}^{+3 .}$ | $0.78{ }_{-0.30}^{+0.38}$ | $6.3_{-2.5}^{+5.7} \times 10^{4}$ | $1.3_{-0.1}^{+0.1} \times 10^{8}$ |
| NGC3265 | 28. ${ }_{-3}^{+5}$. | $1.68{ }_{-0.43}^{+0.38}$ | $6.2_{-3.6}^{+5.5} \times 10^{5}$ | $1.2{ }_{-0.1}^{+0.1} \times 10^{9}$ |
| NGC3351 | 21. ${ }_{-2}^{+3}$. | $2.05{ }_{-0.44}^{+0.43}$ | $9.8{ }_{-5.3}^{+7.9} \times 10^{6}$ | $4.3{ }_{-0.2}^{+0.2} \times 10^{9}$ |
| NGC3521 | $20 .{ }_{-1}^{+2}$. | $2.20_{-0.33}^{+0.32}$ | $6.9_{-2.7}^{+3.9} \times 10^{7}$ | $2.2{ }_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC3621 | 21. ${ }_{-2}^{+3 .}$ | $1.93{ }_{-0.47}^{+0.45}$ | $1.1_{-0.6}^{+1.1} \times 10^{7}$ | $4.7_{-0.2}^{+0.2} \times 10^{9}$ |
| NGC3627 | 20. ${ }_{-0,3}^{+1 .}$ | $2.50_{-0.22}$ | $5.7_{-1.3}^{+0.3} \times 10^{7}$ | $1.7_{-0.1}^{+0.2} \times 10^{10}$ |
| NGC3773 | 23. ${ }_{-3}^{+3 .}$. | $1.922_{-0.38}^{-+0.42}$ | $3.8{ }_{-1.6}^{+1.3} \times 10^{5}$ | $3.0_{-0.2}^{+0.2} \times 10^{8}$ |
| NGC3938 | $20 .{ }_{-2}^{+2}$. | $2.05{ }_{-0.38}^{+0.48}$ | $3.1_{-1.4}^{+2.8} \times 10^{7}$ | $1.0_{-0.04}^{+0.04} \times 10^{10}$ |
| NGC4236 | 26. ${ }_{-4}^{+5 \text {. }}$ | $0.79_{-0.48}^{+0.46}$ | $1.6{ }_{-0.8}^{+1.9} \times 10^{5}$ | $2.99_{-0.1}^{+0.1} \times 10^{8}$ |
| NGC4254 | 20. ${ }_{-0.5}^{+2 .}$ | $2.39_{-0.34}^{+0.09}$ | $7.6_{-3.1}^{+1.3} \times 10^{7}$ | $2.4_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC4321 | 19. ${ }_{-1 .}^{+2 .}$ | $2.48{ }_{-0.43}^{-0.09}$ | $1.1_{-0.4}^{+0.2} \times 10^{8}$ | $1.9_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC4536 | 24. ${ }_{-3}^{+4 .}$ | $1.74_{-0.37}^{+0.49}$ | $1.1_{-0.5}^{+1.2} \times 10^{7}$ | $1.2_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC4559 | 24. ${ }_{-3}^{+4 .}$ | $1.400_{-0.44}^{+0.42}$ | $1.9{ }_{-1.0}^{+2.0} \times 10^{6}$ | $1.88_{-0.1}^{+0.1} \times 10^{9}$ |
| NGC4569 | $20 .{ }_{-1}^{+2}$. | $2.35_{-0.37}^{+0.29}$ | $1.3{ }_{-0.5}^{+0.6} \times 10^{7}$ | $3.3_{-0.1}^{+0.1} \times 10^{9}$ |
| NGC4579 | 18. ${ }_{-1}^{+2}$ | $2.46{ }_{-0.49}^{+0.08}$ | $4.88_{-2.1}^{+1.0} \times 10^{7}$ | $7.2_{-0.3}^{+0.3} \times 10^{9}$ |
| NGC4594 | 19. ${ }_{-3}^{+1 .}$ | $1.97{ }_{-0.36}^{+0.58}$ | $9.5{ }_{-2.8}^{+11.8} \times 10^{6}$ | $2.4{ }_{-0.1}^{+0.1} \times 10^{9}$ |
| NGC4625 | 21. ${ }_{-2}^{+3}$. | $1.76{ }_{-0.39}^{+0.56}$ | $7.3_{-3.5}^{+10.3} \times 10^{5}$ | $3.1_{-0.1}^{+0.1} \times 10^{8}$ |
| NGC4631 | 24. ${ }_{-3}^{+3}$. | $1.83_{-0.42}^{+0.43}$ | $1.6{ }_{-0.7}^{+1.7} \times 10^{7}$ | $1.4{ }_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC4725 | 18. ${ }_{-1}^{+3 .}$ | $2.06_{-0.56}^{+0.51}$ | $3.2{ }_{-1.9}^{+3.0} \times 10^{7}$ | $4.4{ }_{-0.2}^{+0.2} \times 10^{9}$ |
| NGC4736 | 25. | $1.87{ }_{-0.41}^{+0.53}$ | $2.9{ }_{-1.5}^{+2.8} \times 10^{6}$ | $3.6_{-0.2}^{+0.2} \times 10^{9}$ |
| DDO154 | - | - | ${ }_{-}$ | - |
| NGC4826 | $23 .{ }_{-2}^{+3}$. | $2.24_{-0.37}^{+0.32}$ | $4.2{ }_{-1.8}^{+2.4} \times 10^{6}$ | $2.8_{-0.1}^{+0.1} \times 10^{9}$ |
| DDO165 | - | - | - | - |



Fig. 5.12. SEDs of Haro 11 (left) and II Zw 40 (right). They have been obtained with a modified blackbody model with a fixed $\beta_{\text {theo }}=2.0$. The $870 \mu \mathrm{~m}$ point for Haro 11 is from LABOCA (Galametz et al. 2009). For II Zw 40 , the 450 and $850 \mu \mathrm{~m}$ points are from SCUBA, and the 1.2 mm point is from MAMBO (Galliano et al. 2005). All submm points have been corrected for non-dust contamination (free-free and synchrotron radiations and CO line contamination, see Section 6.5). The filled blue circles are the modelled fluxes in each band. The red crosses are the observations. The total SEDs are displayed in black. The bottom panel of each plot indicates the residuals from the fit.

Table 5.2. Table of modified blackbody fit parameters for the KINGFISH galaxies (continued).

| Source | Temperature (K) | $\beta_{\text {obs }}$ | $\mathrm{M}_{\text {dust }, B B}\left(M_{\odot}\right)$ | $\mathrm{L}_{\text {FIR-BB }}\left(L_{\odot}\right)$ |
| :---: | :---: | :---: | :---: | :---: |
| NGC5055 | 19. ${ }_{-1}^{+2 .}$ | $2.32_{-0.42}^{+0.27}$ | $6.22_{-2.4}^{+2.9} \times 10^{7}$ | $1.2{ }_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC5398 | $34 .{ }_{-6 .}$. | $0.588_{-0.36}^{+0.38}$ | $3.1_{-1.6}^{+3.1} \times 10^{4}$ | $2.0_{-0.1}^{+0.1} \times 10^{8}$ |
| NGC5408 | $30 .{ }_{-5}^{+6}$. | $1.67{ }_{-0.44}^{+0.55}$ | $2.1_{-1.1}^{+2.8} \times 10^{4}$ | $1.3{ }_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC5457 | 21. ${ }_{-3}$. | $1.76{ }_{-0.37}^{+0.52}$ | $2.6{ }_{-1.1}^{+3.0} \times 10^{7}$ | $6.5{ }_{-0.4}^{+0.7} \times 10^{7}$ |
| NGC5474 | $23 .{ }_{-2}^{+3}$. | $1.16_{-0.41}^{+0.45}$ | $3.22_{-1.7}^{+3.7} \times 10^{5}$ | $2.7_{-0.1}^{+0.1} \times 10^{8}$ |
| NGC5713 | $22 .{ }_{-1}^{+3}$. | $2.42{ }_{-0.46}^{+0.06}$ | $3.5_{-1.6}^{+0.7} \times 10^{7}$ | $1.9{ }_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC5866 | 21. ${ }_{-1}^{+2}$. | $2.50_{-0.41}^{+0.00}$ | $1.1_{-0.4}^{+0.2} \times 10^{7}$ | $4.1_{-0.2}^{+0.2} \times 10^{9}$ |
| NGC6946 | $20 .{ }_{-1}^{+2}$. | $2.34_{-0.42}^{+0.18}$ | $7.0_{-3.5}^{+1.7} \times 10^{7}$ | $2.2_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC7331 | $20 .{ }_{-1}^{+2}$. | $2.199_{-0.38}^{+0.37}$ | $9.6_{-4.5}^{+5.5} \times 10^{7}$ | $3.1_{-0.1}^{+0.1} \times 10^{10}$ |
| NGC7793 | $23 .{ }_{-2}^{+3}$. | $1.50{ }_{-0.40}^{+0.44}$ | $1.66_{-0.7}^{+1.6} \times 10^{6}$ | $1.2{ }_{-0.05}^{+0.04} \times 10^{9}$ |

Chapter 5. Cold dust from low-metallicity environments to normal galaxies

Table 5.3. Table of relative residuals at $500 \mu \mathrm{~m}$ for a modified blackbody fit with $\beta_{\text {theo }}$ fixed to 2.0 for the DGS and KINGFISH samples. A column with the $\beta_{o b s}$ values from Table 5.1 have been added.

|  | $\beta_{o b s}$ | $\beta_{\text {theo }}=2.0$ |  |  |
| :---: | :---: | :---: | :---: | :---: |
| Source |  | $R(500)$ (\%) | $\Delta R(500)(\%)$ | Excess ? |
| DGS |  |  |  |  |
| Haro11 | 1.96 | 27.6 | 16.7 | yes |
| Haro2 | 2.38 | -6.4 | 11.4 |  |
| Haro3 | 2.15 | -2.3 | 11.0 |  |
| He2-10 | 2.24 | -3.3 | 9.1 |  |
| HS0052+2536 | 1.20 | 154.2 | 125.3 | yes |
| IC10 | 2.25 | 0.1 | 28.4 |  |
| IIZw40 | 1.71 | 2.2 | 8.2 |  |
| Mrk1089 | 2.34 | -3.6 | 13.1 |  |
| Mrk930 | 2.22 | 8.3 | 31.0 |  |
| NGC1140 | 2.17 | -3.6 | 15.8 |  |
| NGC1569 | 2.20 | 1.4 | 9.2 |  |
| NGC1705 | 1.16 | 42.5 | 35.0 | yes |
| NGC2366 | 0.96 | 39.6 | 22.8 | yes |
| NGC4214 | 1.39 | 17.9 | 11.5 | yes |
| NGC4449 | 2.18 | 1.2 | 9.0 |  |
| NGC4861 | 1.36 | 32.5 | 26.6 | yes |
| NGC5253 | 1.86 | 8.5 | 9.8 |  |
| NGC625 | 1.33 | 21.6 | 13.7 | yes |
| NGC6822 | 0.99 | 96.8 | 66.2 | yes |
| UM311 | 1.58 | 9.4 | 10.8 |  |
| UM448 | 1.99 | 6.9 | 18.5 |  |
| VIIZw403 | 1.57 | 71.7 | 58.5 | yes |
| KINGFISH |  |  |  |  |
| NGC0337 | 1.70 | 17.8 | 13.5 | yes |
| NGC0628 | 2.19 | 6.3 | 9.9 |  |
| NGC0855 | 1.41 | 11.8 | 20.6 |  |
| NGC0925 | 1.10 | 35.9 | 15.3 | yes |
| NGC1097 | 1.98 | 7.6 | 11.3 |  |
| NGC1266 | 2.31 | -1.9 | 12.6 |  |
| NGC1291 | 1.70 | 10.8 | 12.4 |  |
| NGC1316 | 2.50 | 2.7 | 18.8 |  |
| NGC1377 | 1.79 | 26.1 | 21.2 | yes |
| IC0342 | 2.17 | 2.1 | 9.1 |  |
| NGC1482 | 2.36 | -1.5 | 10.4 |  |
| NGC1512 | 0.98 | 32.3 | 15.3 | yes |
| NGC2146 | 2.50 | -6.8 | 8.2 |  |
| HoII | 1.88 | 26.4 | 72.3 |  |
| NGC2798 | 2.24 | 4.6 | 10.3 |  |
| NGC2841 | 2.33 | 1.4 | 9.4 |  |
| NGC2915 | 0.73 | 64.2 | 25.8 | yes |
| HoI | 0.00 | 98.1 | 73.9 | yes |
| NGC2976 | 1.99 | 7.9 | 11.1 |  |
| NGC3049 | 0.82 | 63.1 | 18.4 | yes |
| NGC3077 | 1.62 | 22.1 | 12.8 | yes |
| M81dwB | 0.82 | 62.0 | 87.8 |  |

Table 5.3. (continued).

|  | $\beta_{\text {obs }}$ |  |  |  |
| :--- | :---: | :---: | :---: | :---: |
| Source |  | $R(500)(\%)$ | $\Delta R(500)(\%)$ | Excess ? |
| NGC3190 | 2.50 | -5.1 | 9.2 |  |
| NGC3184 | 2.18 | 5.3 | 11.0 |  |
| NGC3198 | 1.58 | 18.4 | 10.9 | yes |
| IC2574 | 0.78 | 38.7 | 17.9 | yes |
| NGC3265 | 1.68 | 29.5 | 22.0 | yes |
| NGC3351 | 2.05 | 2.8 | 8.5 |  |
| NGC3521 | 2.20 | 3.0 | 9.9 |  |
| NGC3621 | 1.93 | 13.5 | 11.5 | yes |
| NGC3627 | 2.50 | -10.8 | 8.9 |  |
| NGC3773 | 1.92 | 13.0 | 17.2 |  |
| NGC3938 | 2.05 | 6.3 | 11.9 |  |
| NGC4236 | 0.79 | 54.9 | 19.1 | yes |
| NGC4254 | 2.39 | -6.7 | 9.0 |  |
| NGC4321 | 2.48 | -2.1 | 9.8 |  |
| NGC4536 | 1.74 | 17.1 | 12.1 | yes |
| NGC4559 | 1.40 | 26.7 | 14.0 | yes |
| NGC4569 | 2.35 | 0.1 | 9.0 |  |
| NGC4579 | 2.46 | -1.1 | 9.0 |  |
| NGC4594 | 1.97 | 9.9 | 9.0 | yes |
| NGC4625 | 1.76 | 16.8 | 14.9 | yes |
| NGC4631 | 1.83 | 15.5 | 12.7 | yes |
| NGC4725 | 2.06 | 9.0 | 10.6 |  |
| NGC4736 | 1.87 | 13.9 | 11.5 | yes |
| NGC4826 | 2.24 | 2.9 | 9.7 |  |
| NGC5055 | 2.32 | 1.3 | 9.8 |  |
| NGC5398 | 0.58 | 62.0 | 22.7 | yes |
| NGC5408 | 1.67 | 79.3 | 40.4 | yes |
| NGC5457 | 1.76 | 14.8 | 12.9 | yes |
| NGC5474 | 1.16 | 36.4 | 15.5 | yes |
| NGC5713 | 2.42 | -1.8 | 9.8 |  |
| NGC5866 | 2.50 | -4.1 | 9.7 |  |
| NGC6946 | 2.34 | 0.1 | 8.9 |  |
| NGC7331 | 2.19 | 3.5 | 9.3 |  |
| NGC7793 | 1.50 | 23.4 | 12.7 | yes |
|  |  |  |  |  |



Fig. 5.13. Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the DGS. The solid black line is the modelled modified blackbody, the blue circles are the modelled points. The red symbols are the observations: crosses are for detections and diamonds are for upper limits. We overlaid the MIPS $24 \mu \mathrm{~m}$ point from Bend 9 et al. (2012). The $T$ and $\beta$ parameters have been indicated on the top of each plot along with the $\chi^{2}$ value of the fit. The bottom panel of each plot indicates the residuals from the fit.


Fig. 5.13. (continued) Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the DGS.


Fig. 5.13. (continued) Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the DGS. Note: for SBS0335052 , we included the $24 \mu \mathrm{~m}$ point in the fit as the $24 \mu \mathrm{~m}$ point fell below the modelled modified blackbody when we just overlaid it on the plot.


Fig. 5.13. (continued) Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the DGS.


Fig. 5.13. (continued) Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the DGS. Note: for SBS1211 +540 , we included the $24 \mu \mathrm{~m}$ point in the fit as the $24 \mu \mathrm{~m}$ point fell below the modelled modified blackbody when we just overlaid it on the plot.


Fig. 5.14. Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the KINGFISH galaxies. The solid black line is the modelled modified blackbody, the blue circles are the modelled points. The red symbols are the observations: crosses are for detections and diamonds are for upper limits. We overlaid the MIPS 24129 point from Dale et al. (2007). The $T$ and $\beta$ parameters have been indicated on the top of each plot along with the $\chi^{2}$ value of the fit. The bottom panel of each plot indicates the residuals from the fit.


Fig. 5.14. (continued) Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the KINGFISH galaxies.


Fig. 5.14. (continued) Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the KINGFISH galaxies.


Fig. 5.14. (continued) Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the KINGFISH galaxies.


Fig. 5.14. (continued) Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the KINGFISH galaxies.


Fig. 5.14. (continued) Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the KINGFISH galaxies.


Fig. 5.14. (continued) Modified blackbody fits of the 70 to $500 \mu \mathrm{~m}$ range for the KINGFISH galaxies.

## Part III

## Evolution of ISM properties with metallicity : a complete IR-to-submillimetre view

## Chapter 6

## Sampling the IR-to-submm observed dust SEDs

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In this Chapter we collect all of the ancillary data needed to properly sample the IR-to-submm dust SEDs of the DGS galaxies. We intensively use the archives whenever it is possible. Otherwise we treat the data ourselves. References for finding the corresponding KINGFISH data we use is given at the end of each section.

### 6.1 2MASS: Searching the literature

2MASS (Two Microns All Sky Survey) observed the entire sky in the J $(1.24 \mu \mathrm{~m})$, H $(1.66 \mu \mathrm{~m})$ and $\mathrm{K}_{s}(2.16 \mu \mathrm{~m})$ bands. The RSRFs (relative spectral response functions) for each band is shown in Fig. 6.1. The RSRF for 2MASS are designed to be integrated directly over spectra in $\mathrm{F}_{\lambda}$ (with $\nu F_{\nu}=\lambda F_{\lambda}$ ), in order to calculate synthetic photometry.

The 2MASS data for the DGS sample is compiled from the literature: the NASA/IPAC ISA 2MASS Point Source Catalog, the 2MASS Extended Objects Final Release, the 2MASS Large Galaxy Atlas (Jarrett et al. 2003), Engelbracht et al. (2008); Dale et al. (2009) and is given in Table 6.8. In some cases, the original data is given in magnitudes in the databases. To convert this into flux densities, we use the zero-magnitude flux values from Cohen et al. (2003), $F_{0}$ :

$$
\begin{equation*}
F_{\nu}=F_{0} \times 10^{-m a g / 2.5} \tag{6.1}
\end{equation*}
$$

The error on the flux density is given by Eq. 6.2 and is derived by propagating the errors on $F_{0}$ and on the magnitude, mag, $\sigma_{F_{0}}$ and $\sigma_{m a g}$, in Eq. 6.1. $F_{0}$ and $\sigma_{F_{0}}$ are given in Table 6.1 along with some characteristics of the 2MASS data.

$$
\begin{equation*}
\sigma_{F_{\nu}}=F_{\nu} \times \sqrt{\left(\frac{\sigma_{F_{0}}}{F_{0}}\right)^{2}+\left(\frac{\ln (10)}{2.5} \sigma_{m a g}\right)^{2}} \tag{6.2}
\end{equation*}
$$

Table 6.1. Characteristics of 2MASS (from Cohen et al. 2003; Jarrett et al. 2003).

| Band | $\lambda[\mu \mathrm{m}]$ | $F_{0}[\mathrm{Jy}]$ | $\sigma_{F_{0}}[\mathrm{Jy}]$ | Calibration accuracy |
| :--- | :---: | :---: | :---: | :---: |
| J | 1.25 | 1594.0 | 27.8 | $3 \%$ |
| H | 1.65 | 1024.0 | 20.0 | $3 \%$ |
| Ks | 2.20 | 666.7 | 12.6 | $3 \%$ |

KINGFISH 2MASS photometry is presented in Dale et al. (2007) and has been taken from Jarrett et al. (2003).

### 6.2 Spitzer observations

Spitzer observations of some dwarf galaxies of the DGS were already available through the Spitzer database ${ }^{1}$. When the DGS was designed, complementary Spitzer observations were scheduled in the cycle 5 program to complete the set of existing Spitzer data in order to have NIR and MIR complementary observations for a maximum number of targets in the DGS sample (Dust Evolution in Low-Metallicity Environments: P.I. F. Galliano; ID: 50550). These new observations are presented here, along with those previously available.

[^24]

Fig. 6.1. 2MASS filter transmissions, $R(\lambda)$. The dashed lines mark the reference wavelength for each band.

### 6.2.1 IRAC

## Instrument

The InfraRed Array Camera (IRAC) is a four-channel camera providing simultaneous $5.2^{\prime} \times 5.2^{\prime}$ images at $3.6,4.5,5.8$ and $8.0 \mu \mathrm{~m}$. All four detectors arrays are $256 \times 256$ pixels in size, with pixel sizes of $1.2^{\prime \prime}$. The main characteristics of IRAC are given in Table 6.2. Several stars were used for the calibration. The total calibration uncertainty is $\sim 10 \%$.

The convention for the spectral response function here is that $\mathrm{R}(\nu)$ is proportional to the number of electrons produced by a single photon of energy $h \nu$. Thus the number of electrons collected from the source, $N_{e}^{S}$, during an integration time $t$ and though a telescope area $\mathcal{A}$, is directly proportional to the source flux density, $F_{S}(\nu)$, divided by the energy of the collected photons, weighted by the spectral response of the telescope, $\mathrm{R}(\nu)$ :

$$
\begin{equation*}
N_{e}^{S}=t \mathcal{A} \int \frac{F_{S}(\nu)}{h \nu} R(\nu) \mathrm{d} \nu \tag{6.3}
\end{equation*}
$$

The spectral response for each band is shown in Fig. 6.2. Here the spectral convention is $\nu \times F_{\nu}=\operatorname{cst}$ or $\mathrm{F}_{S}(\nu)=\mathrm{F}_{S}\left(\nu_{0}\right) \times\left(\nu_{0} / \nu\right)$, with $\nu_{0}$ corresponding to $\lambda_{0}=\{3.6,4.5,5.8,8.0\} \mu \mathrm{m}$. Importing this into Eq. 6.3 gives:

$$
\begin{equation*}
N_{e}^{S}=\frac{t \mathcal{A} F_{S}\left(\nu_{0}\right)}{h \nu_{0}} \int\left(\frac{\nu_{0}}{\nu}\right)^{2} R(\nu) \mathrm{d} \nu \tag{6.4}
\end{equation*}
$$

and thus:

$$
\begin{equation*}
F_{S}\left(\nu_{0}\right)=\frac{t \mathcal{A}}{N_{e}^{S} h \nu_{0}} \int\left(\frac{\nu_{0}}{\nu}\right)^{2} R(\nu) \mathrm{d} \nu \tag{6.5}
\end{equation*}
$$

However, as for PACS and SPIRE, there is no particular reason for the source to have a spectral shape $\mathrm{F}_{S}(\nu) \propto \nu^{-1}$. Thus we need to apply a colour correction. If the source "true" spectral shape, $\mathrm{F}_{S}^{\prime}(\nu)$, is unknown, we have from Eq. 6.3:

$$
\begin{equation*}
N_{e}^{\prime S}=t \mathcal{A} \int \frac{F_{S}^{\prime}(\nu)}{h \nu} R(\nu) \mathrm{d} \nu \tag{6.6}
\end{equation*}
$$

By definition the calibration factor, $C$, is the ratio between the number of electrons collected from an arbitrary source to the flux density of this source at the nominal wavelength, i.e., $\mathrm{N}_{e} / \mathrm{F}\left(\nu_{0}\right)$. Using the fact that: $\mathrm{N}_{e}^{S} / \mathrm{F}_{S}\left(\nu_{0}\right)=\mathrm{N}_{e}^{\prime S} / \mathrm{F}_{S}^{\prime}\left(\nu_{0}\right)$, we get:

$$
\begin{equation*}
F_{S}^{\prime}\left(\nu_{0}\right)=t \mathcal{A} \int \frac{F_{S}^{\prime}(\nu)}{h \nu} R(\nu) \mathrm{d} \nu \times \frac{h \nu_{0}}{t \mathcal{A} \int\left(\nu_{0} / \nu\right)^{2} R(\nu) \mathrm{d} \nu} \tag{6.7}
\end{equation*}
$$

which gives in the end:

$$
\begin{equation*}
F_{S}^{\prime}\left(\nu_{0}\right)=\frac{\int F_{S}^{\prime}(\nu)\left(\nu_{0} / \nu\right) R(\nu) \mathrm{d} \nu}{\int\left(\nu_{0} / \nu\right)^{2} R(\nu) \mathrm{d} \nu} \tag{6.8}
\end{equation*}
$$

Eq. 6.8 can be generalised to any flux convention in the form $\mathrm{F}_{S}(\nu) \propto \nu^{-\alpha_{S}}$ chosen for a photomultiplier (such as MIPS, WISE or IRAS presented in the next Sections):

$$
\begin{equation*}
F_{S}^{\prime}\left(\nu_{0}\right)=\frac{\int F_{S}^{\prime}(\nu)\left(\nu_{0} / \nu\right) R(\nu) \mathrm{d} \nu}{\int\left(\nu_{0} / \nu\right)^{\alpha_{S}+1} R(\nu) \mathrm{d} \nu} \tag{6.9}
\end{equation*}
$$



Fig. 6.2. IRAC filter transmissions, $R(\lambda)$. The dashed lines mark the reference wavelength for each band.

Table 6.2. IRAC characteristics (from the IRAC Instrument Handbook, version 2.0.1).

| IRAC |  |  |  |  |  |  |  |  |
| :--- | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Band | Wavelength $(\mu \mathrm{m})$ | Bandwidth $(\mu \mathrm{m})$ | FWHM $\left({ }^{\prime \prime}\right)$ | Calibration accuracy | $\mathrm{A}^{a}$ | $\mathrm{~B}^{a}$ | $\mathrm{C}^{a}$ |  |
| IRAC 1 | 3.6 | 0.750 | 1.66 | $10 \%$ | 0.82 | 0.370 | 0.910 |  |
| IRAC 2 | 4.5 | 1.015 | 1.72 | $10 \%$ | 1.16 | 0.433 | 0.94 |  |
| IRAC 3 | 5.8 | 1.425 | 1.88 | $10 \%$ | 1.49 | 0.207 | 0.66 |  |
| IRAC 4 | 8.0 | 2.905 | 1.98 | $10 \%$ | 1.37 | 0.330 | 0.740 |  |

[^25]
## Photometry

The IRAC maps are retrieved from the NASA/IPAC ISA database for Spitzer data, and the AORkeys for the IRAC observations are summarised in Appendix A. After a step of background subtraction, similar to what was done for SPIRE (Chapter 4), flux densities are extracted from the maps through aperture photometry, and are given in Table 6.9. The apertures are in most cases the same as the ones used for Herschel photometry. Some contamination which can occur from background sources or foreground stars has also been removed from the source aperture.

The IRAC calibration is based on point-source photometry for a $12^{\prime \prime}$ radius aperture. Extra aperture correction is needed to account for the emission from the wings of the PSF and the scattering of diffuse emission across the IRAC focal plane. This correction is given by:

$$
\begin{equation*}
\frac{F_{\nu_{0}}(\text { true })}{F_{\nu_{0}}(\text { measured })}=A\left(\nu_{0}\right) e^{-r^{B\left(\nu_{0}\right)}}+C\left(\nu_{0}\right) \tag{6.10}
\end{equation*}
$$

where, $A\left(\nu_{0}\right), B\left(\nu_{0}\right)$ and $C\left(\nu_{0}\right)$ are given in Table 6.2 and $r$ is the source aperture radius. Following the recipe given in the IRAC Instrument Handbook (version 2.0.1 Section 4.11.1 ${ }^{2}$ ) we do not apply this correction to small and compact sources (see Table 6.9).

## Errors

The uncertainty on the flux density is determined the same way as for SPIRE. Adapting Eq. 4.9 to IRAC gives:

$$
\begin{equation*}
u n c_{f l u x}=\sqrt{u n c_{b g}^{2}+u n c_{\text {source }}^{2}} \tag{6.11}
\end{equation*}
$$

where:

$$
\left\{\begin{align*}
u n c_{b g} & =N_{a p} \frac{\sigma_{s k y}}{\sqrt{N_{b g}}}  \tag{6.12}\\
\text { unc }_{\text {source }} & =\sqrt{N_{a p} \sigma_{s k y}}
\end{align*}\right.
$$

where $\sigma_{s k y}$ is the standard deviation of all of the pixels in the background aperture. $N_{a p}$ and $N_{b g}$ are the number of pixels in the source and background apertures respectively. The final total uncertainty, $\sigma_{3.6-4.5-5.8-8.0}$ reported in Table 6.9, is obtained by adding $u n c_{f l u x}$ and the $10 \%$ calibration error in quadrature.

We consider that galaxies are not detected when the computed flux density is lower than three times the corresponding uncertainty on the flux density in a given band. The final upper limit is then three times the uncertainty on the flux density value ( $3 \sigma$ upper limit, reported in Table 6.9).

IRAC photometry for KINGFISH galaxies is presented in Dale et al. (2007).

## Comparison with the literature

IRAC photometry is available in the literature for 29 DGS sources: Hunt et al. (2006); Dale et al. (2007, 2009); Engelbracht et al. (2008); Galametz et al. (2009). We use the ratios of our IRAC to the literature IRAC flux densities to assess how well the measurements agree with each other; a ratio of 1 corresponds to a very good agreement. If we compare our measurements to those from the literature we get a fairly good agreement between them (see Fig. 6.3), with some outliers:

Haro2: The comparison to the literature values agrees for the first three IRAC bands (ratios of $0.92,0.96$ and 0.93 at $3.6 \mu \mathrm{~m}, 4.5 \mu \mathrm{~m}$, and $5.8 \mu \mathrm{~m}$ respectively) except for the IRAC $8.0 \mu \mathrm{~m}$ band

[^26]where the ratio goes to $\sim 2$. We check the photometry at $8.0 \mu \mathrm{~m}$ by recomputing the flux density with the aperture given in Dale et al. (2007) and find a flux in very close agreement to the one we find with our own aperture (ratio of 1.009 at $8.0 \mu \mathrm{~m}$ ), and also for the three other IRAC bands. As an additional check, we find that Marble et al. (2010) performed photometry for Haro2 in a IRSmatched aperture and their result is in good agreement with ours (ratio $=1.036$ at $8.0 \mu \mathrm{~m}$ ). Thus we discard the $8.0 \mu \mathrm{~m}$ point from Dale et al. (2007) in our comparison with literature measurements.

NGC4861: Our measurements are systematically larger than the flux from Engelbracht et al. (2008). A possible explanation could be that we include Mrk39 with which NGC4861 is interacting in our aperture and that Engelbracht et al. (2008) did not. However, no precise information is given in Engelbracht et al. (2008) about the size of the apertures used for IRAC photometry, so it is hard to see if we are really comparing similar measurements. Thus we discard this galaxy in our comparison with literature measurements.


Fig. 6.3. Comparison of our IRAC flux densities and literature IRAC flux densities: $\mathrm{F}_{\text {IRAC }}\left(\lambda_{0}\right) / \mathrm{F}_{\text {IRACLIT }}\left(\lambda_{0}\right)$ flux density ratios as a function of our IRAC flux density, $\mathrm{F}_{\text {IRAC }}\left(\lambda_{0}\right)$, at 3.6 $\mu \mathrm{m}$ (top left), $4.5 \mu \mathrm{~m}$ (top right), $5.8 \mu \mathrm{~m}$ (bottom left) and $8.0 \mu \mathrm{~m}$ (bottom right). As a guide to the eye, the unity line is added as a solid line as well as the average uncertainties on the ratio in each bands as dotted lines. These average uncertainties are $17 \%, 17 \%, 19 \%$ and $17 \%$ at $3.6,4.5,5.8$ and $8.0 \mu \mathrm{~m}$. Colours distinguish the reference for the literature measurement.

UGC4483: We do not detect this galaxy at $5.8 \mu \mathrm{~m}$. For the other three IRAC bands, our measurements are systematically larger than the flux density from Engelbracht et al. (2008). However, no precise information is given in Engelbracht et al. (2008) about the size of the aperture used for IRAC photometry. Nonetheless they mention an aperture for 2MASS measurements which is very similar to the one we use. IRAC photometry within this 2MASS aperture is consistent with what we find with our aperture. Thus we discard this galaxy in our comparison with literature measurements.

Tol1214-27\%: The comparison is fine at all wavelengths except for IRAC $5.8 \mu \mathrm{~m}$. Again, because of the lack of information about the apertures used in Engelbracht et al. (2008) we cannot assess that we are really comparing similar measurements. Nonetheless, the ratio between the two measurements is coherent, within its error bar, with photometric agreement because of the large uncertainties on both measurements ( $\sim 20 \%$ (Table 6.9) and $41 \%$ (Engelbracht et al. 2008)). Thus we keep this point in our comparison with literature measurements.

Excluding these outliers, and points where one of the two measurements (ours or the literature's) is an upper limit, we get a mean ratio of our IRAC measurements to those in the literature of 1.05 $\pm 0.11$ for IRAC $3.6 \mu \mathrm{~m}, 1.06 \pm 0.11$ for IRAC $4.5 \mu \mathrm{~m}, 1.02 \pm 0.14$ for IRAC $5.8 \mu \mathrm{~m}$ and 1.01 $\pm 0.11$ for IRAC $8.0 \mu \mathrm{~m}$. This is to be compared to an average uncertainty on the ratios of $17 \%$, $17 \%, 19 \%$ and $17 \%$ at $3.6,4.5,5.8$ and $8.0 \mu \mathrm{~m}$. The error on the average ratio is comparable to the average uncertainties on the ratio for all of the bands. Thus there is a good photometric agreement between our measurements and previously published measurements.

### 6.2.2 IRS

## Instrument

The InfraRed Spectrograph ${ }^{3}$ (IRS) is composed of four modules providing low ( $\mathrm{R} \sim 60-130$ ) and moderate ( $\mathrm{R} \sim 600$ ) resolution spectroscopic capabilities over 5.8-40 $\mu \mathrm{m}$ (ShortLow and LongLow modules) and over 9.9-37.2 $\mu \mathrm{m}$ (ShortHigh and LongHigh modules). IRS also provide imaging in two filters: Blue ( $13-18 \mu \mathrm{~m}$ ) and Red ( $18-26 \mu \mathrm{~m}$ ). The basic informations for the four spectroscopy modules can be found in Table 6.3, and a schematic view of the IRS slits is shown in Fig. 6.4.

Table 6.3. IRS characteristics (from the IRS Instrument Handbook, version 5.0).

| IRS |  |  |  |  |  |  |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Module | Range ( $\mu \mathrm{m}$ ) | Resolving Power R | Plate scale ("/pix) | Slit Width (') | Slit Length (") | Cut-offs ( $\mu \mathrm{m}$ ) |
| SL1 | 7.46-14.29 | 61-120 |  | 3.6 |  | [ $\lambda_{\text {min }}-7.53$, |
| SL2 SL3 | $5.13-7.60$ $7.33-8.66$ | 60-127 | 1.8 | 3.7 | 57 | $\begin{aligned} & \left.14.02-\lambda_{\max }\right] \\ & {\left[\lambda_{\text {min }}-5.23,\right.} \\ & \left.7.49-\lambda_{\max }\right] \\ & {\left[\lambda_{\text {min }}-7.4,\right.} \\ & \left.8.5-\lambda_{\text {max }}\right] \end{aligned}$ |
| LL1 | 19.91-39.90 | 58-112 |  | 10.5 |  | $\begin{aligned} & {\left[\lambda_{\text {min }}-20.5,\right.} \\ & \left.37.4-\lambda_{\text {max }}\right] \end{aligned}$ |
| LL2 LL3 | $13.90-21.27$ $19.23-21.61$ | 57-126 | 5.1 | 10.7 | 168 | $\begin{aligned} & {\left[\lambda_{\text {min }}-14.0,\right.} \\ & \left.20.52-\lambda_{\text {max }}\right] \\ & {\left[\lambda_{\text {min }}-19.8,\right.} \\ & \left.21.5-\lambda_{\text {max }}\right] \end{aligned}$ |
| SH | 9.89-19.51 | 600 | 2.3 | 4.7 | 11.3 |  |
| LH | 18.83-37.14 | 600 | 4.5 | 11.1 | 22.3 |  |

[^27]

Fig. 6.4. Schematic view of the IRS slits. Note that the slits are not parallel in the Spitzer plane as shown here. Actual slit position angles relative to a Spitzer roll angle of $0^{\circ}$ are $\mathrm{SL}=+84.7^{\circ}, \mathrm{LL}=+181.2^{\circ}, \mathrm{SH}=$ $+221.5^{\circ}, \mathrm{LH}=+136.7^{\circ}$, and IRS Peak-up $=+177.0^{\circ}$.

## Reduction

The IRS spectra for the DGS galaxies have been extracted from the Cornell AtlaS of Spitzer Infrared spectrograph Sources ${ }^{4}$ (CASSIS v5, Lebouteiller et al. 2011) which provides optimal extraction of a large collection of spectra from staring mode observations. Most of the DGS galaxies were observed using the staring mode, except for a few extended ones that were observed with the mapping mode, for which the spectra were reduced manually. We were able to obtain SL and LL data for 43 galaxies in total.

Two versions of the spectra are available in CASSIS: an "optimal" extraction better suited for point sources and a "tapered column" extraction, better suited for partially extended sources. A message advises the user if the "tapered column" extraction or the "optimal" extraction method should be chosen. The chosen extraction for each of the DGS targets, as well as AORkeys, are summarised in Appendix A. CASSIS uses the AdOpt algorithm (Advanced Optimal extraction, Lebouteiller et al. 2010) within the Spectroscopic Modeling Analysis and Reduction Tool (SMART, Higdon et al. 2004). AdOpt enables optimal extraction of spectra using a super-sampled PSF. The pipeline includes such steps as image cleaning, individual exposure combination and background subtraction. Specific attention is given to identification and removal of bad pixels and outlier rejection at the image and spectra levels.

## Preparing the spectra

The IRS spectra for the DGS galaxies are presented in Appendix ??. Several steps are necessary before including them in the SEDs. First some wavelengths at the edges of each orders have to be

[^28]cut off the spectra. These cut-off wavelengths are given in Table 6.3.
For the extended galaxies, there is a conversion step from MJy/sr to Jy where the IRS spectrum is multiplied by the area over which it has been extracted.

We collect IRS data in order to constrain the MIR slope of the SED, thus we are more interested in the continuum rather than in the spectral lines. It is then necessary to rescale the SL and LL spectra in order to match the photometry. The idea is to derive synthetic IRS photometry to correct the two modules and to use all of the constraints we have in this wavelength range: the IRAC 5.8 and $8.0 \mu \mathrm{~m}$, WISE 12 and $22 \mu \mathrm{~m}$ and MIPS $24 \mu \mathrm{~m}$ bands. IRAC 5.8, 8.0 and WISE 12 $\mu \mathrm{m}$ are used simultaneously to derive a correction factor for the SL module that depends on the wavelength. However, for the LL module, we only have two constraints that do not sample well the LL spectrum. Deriving a wavelength-dependant correction factor is not possible here, and the LL correction factor is thus a constant. We use MIPS $24 \mu \mathrm{~m}$ for most cases, while WISE $22 \mu \mathrm{~m}$ is only considered whenever MIPS $24 \mu \mathrm{~m}$ is not available. SL and LL can be treated separately as they are two independent observations.

The correction factor for $\mathrm{LL}, C_{L L}$ is then:

$$
\begin{equation*}
C_{L L}=F_{24} / F_{I R S}\left(M_{24}\right) \tag{6.13}
\end{equation*}
$$

where $F_{24}$ is the observed $24 \mu \mathrm{~m}$ MIPS flux density (from Bendo et al. (2012)) and $F_{I R S}\left(M_{24}\right)$ is the synthetic photometry for the IRS LL, spectrum at $24 \mu \mathrm{~m} . F_{I R S}\left(M_{24}\right)$ is derived using Eq. 6.16. Eq. 6.13 is adapted for WISE $22 \mu \mathrm{~m}$ when MIPS $24 \mu \mathrm{~m}$ is not available.

As we have several constraints available for the SL module, we can use them simultaneously to derive a correction that smoothly depends on wavelength, $C_{S L}(\lambda)$. To derive $C_{S L}(\lambda)$, we fit a spline function, using the IDL spline_P procedure, to $C_{S L}(I 3), C_{S L}($ I4 $), C_{S L}(W 3)$ and $C_{L L}$. We use $C_{L L}$ here to have a better constraint at the end of the spline. We also impose a derivative of 0 at the end of the spline in order to have a smooth function. We also assume that the correction factor $C_{S L}(\lambda)$ should always be $\geq 1.0$.
$C_{S L}(I 3), C_{S L}\left(I_{4}\right), C_{S L}(W 3)$ are given by:

$$
\left\{\begin{align*}
C_{S L}(I 3) & =F_{5.8} / F_{I R S}\left(I_{5.8}\right)  \tag{6.14}\\
C_{S L}(I 4) & =F_{8.0} / F_{I R S}\left(I_{8.0}\right) \\
C_{S L}(W 3) & =F_{12} / F_{I R S, S L}\left(W_{12}\right)
\end{align*}\right.
$$

where $F_{5.8}, F_{8.0}$ and $F_{12}$ are the observed 5.8 and $8.0 \mu \mathrm{~m}$ IRAC and the $12 \mu \mathrm{~m}$ WISE flux densities (from Tables 6.9 and 6.10 ); and $F_{I R S}\left(I_{5.8}\right), F_{I R S}\left(I_{8.0}\right)$ and $F_{I R S, S L}\left(W_{12}\right)$ are the synthetic photometry for the IRS SL spectrum at $5.8,8.0$ and $12 \mu \mathrm{~m} . F_{I R S}\left(I_{5.8}\right), F_{I R S}\left(I_{8.0}\right)$ are derived using Eq. 6.8.

However, there is an overlap between the WISE $12 \mu \mathrm{~m}$ filter and the LL wavelengths (Fig. 6.5). As the LL part of the spectrum has already been corrected, we must find the flux missing in the SL part of the spectrum, i.e., $F_{I R S, S L}\left(W_{12}\right)$, to match $F_{12}$. But the integration over the WISE bandpass is not linear, i.e.,: $F_{I R S, S L}\left(W_{12}\right) \neq F_{I R S}\left(W_{12}\right)-F_{I R S, L L}\left(W_{12}\right)$. In other words, we can not use Eq. 6.18 on the SL wavelength range to derive $F_{I R S, S L}\left(W_{12}\right)$. Thus there is no simple way of deriving $C_{S L}(W 3)$. Instead, we apply the following method:

1. We generate a grid of potential $C_{S L}(W 3),\{\mathrm{x}\}$, from 0.01 to 10 ;
2. find the spline going through $C_{S L}(I 3), C_{S L}\left(I_{4}\right), C_{L L}$, and each $\mathrm{x}_{i}, \mathrm{sp}\left(\mathrm{x}_{i}\right)$;
3. correct the SL spectrum with each $\mathrm{sp}\left(\mathrm{x}_{i}\right)$ and
4. compute the synthetic WISE $12 \mu \mathrm{~m}$ photometry for each corrected total IRS spectrum, $F_{I R S}\left(W_{12}\right)\left(\operatorname{sp}\left(x_{i}\right)\right)$.


Fig. 6.5. (left) Illustration of the overlap between the WISE $12 \mu \mathrm{~m}$ band and the IRS LL wavelengths. (right) Example of a spline used to rescale SL in the case of HS1304+3529. The spline has been fitted to the blue diamonds, except for the $5.8 \mu \mathrm{~m}$ band as the galaxy is not detected at this wavelength. The diamonds represent the ratio between the observed and the synthetic photometry at each wavelength. The black curve is the final adopted spline, $\mathrm{sp}\left(\mathrm{x}_{0}\right)$, where $x_{0}=1.3$. The orange crosses are the corrections applied to the SL part of the spectrum.
$C_{S L}(W 3)$ is the $\mathrm{x}_{0}$ that gives $F_{I R S}\left(W_{12}\right)\left(s p\left(x_{0}\right)\right)=F_{W_{12}}$.
The final IRS spectrum, $F_{I R S, c o r r}\left(\lambda_{S L}\right)$ and $F_{I R S, c o r r}\left(\lambda_{L L}\right)$, is given by:

$$
\left\{\begin{array}{l}
F_{I R S, \text { corr }}\left(\lambda_{S L}\right)=F_{I R S}\left(\lambda_{S L}\right) \times C_{S L}\left(\lambda_{S L}\right)  \tag{6.15}\\
F_{I R S, c o r r}\left(\lambda_{L L}\right)=F_{I R S}\left(\lambda_{L L}\right) \times C_{L L}
\end{array}\right.
$$

The method presented here is adapted depending on the number of constraints for each galaxy. Upper limits are not considered for the correction of IRS spectrum. An example of this treatment of IRS spectrum is shown in Fig. 6.6 along with the spline used to correct the SL spectrum in Fig. 6.5. Note that this correction assumes that the spectral shape of the area observed by the IRS slits likewise describes the expected spectral shape of the full galaxy. This is true for compact sources but can be erroneous for more extended galaxies, except if the region falling within the IRS slits dominates the total emission of the galaxy in the MIR.

For some galaxies, the IRS spectrum is very noisy (e.g. HS1304+3529, Fig 6.6, top panel). To improve the quality of the spectrum, we smooth the spectrum until we reach a signal-to-noise of 3 for each point in the spectrum. This smoothing step is applied for 22 DGS galaxies.

### 6.2.3 MIPS

## Instrument

The Multiband Imaging Photometer for Spitzer (MIPS) provides broad-band imaging and photometry at 24,70 and $160 \mu \mathrm{~m}$. The broad-band filters are shown in Fig. 6.7. The three detector arrays are separated and view the sky simultaneously. The size of the arrays, FOV and pixel sizes are given in Table 6.4. Primary and secondary calibrators for MIPS are stars. MIPS colours corrections are based on a $10^{4} \mathrm{~K}$ blackbody, i.e., $\mathrm{F}_{S}(\nu) \propto B_{\nu}\left(\nu_{0}, T_{0}=10^{4} \mathrm{~K}\right)$ where $\mathrm{B}_{\nu}\left(\nu, T_{0}\right)$ is the Planck function. Eq. 6.9 can thus be transformed into:


Fig. 6.6. Example of IRS treatment for HS1304+3529: the raw spectrum is presented in the top panel. The photometry points have been overlaid in filled purple diamonds as well as synthetic photometry as blue open diamonds. The galaxy is not detected at $5.8 \mu \mathrm{~m}$ as indicated by the open purple triangle. Green, orange and red parts of the spectrum represents the SL2 and LL2, SL3 and LL3, SL1 and LL1 spectra, respectively. The middle panel shows the spectra after cutting-off the bad wavelengths and rescaling to match the photometry. The spline used to rescale SL is shown in Fig. 6.5. The bottom panel shows the spectra smoothed to a S/N of 3 .

$$
\begin{equation*}
F_{S}^{\prime}\left(\nu_{0}\right)=\frac{\int F_{S}^{\prime}(\nu)\left(\nu_{0} / \nu\right) R(\nu) \mathrm{d} \nu}{\int\left(\frac{\nu_{0}}{\nu}\right)\left(\frac{B_{\nu}\left(\nu, T_{0}\right)}{B_{\nu}\left(\nu_{0}, T_{0}\right)}\right) R(\nu) \mathrm{d} \nu} \tag{6.16}
\end{equation*}
$$

with $\mathrm{T}_{0}=10^{4} \mathrm{~K}$.
Table 6.4. MIPS characteristics (from the MIPS Instrument Handbook, version 2).

| MIPS |  |  |  |  |  |  |  |  |
| :--- | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Band | $\lambda$ <br> $(\mu \mathrm{m})$ | Bandwidth <br> $(\mu \mathrm{m})$ | FWHM <br> $\left({ }^{\prime \prime}\right)$ | Calibration accuracy | FOV <br> $\left({ }^{\prime} \times{ }^{\prime}\right)$ | Array size <br> $($ pix $\times$ pix $)$ | Pixel size <br> $\left({ }^{\prime \prime} \times{ }^{\prime \prime}\right)$ |  |
| MIPS 1 | 24 | 4.7 | 6 | $4 \%$ | $5.4 \times 5.4$ | $128 \times 128$ | $2.5 \times 2.6$ |  |
| MIPS 2 | 70 | 19 | 18 | $7 \%$ | $5.2 \times 2.6$ | $32 \times 16$ | $9.9 \times 10.1$ |  |
| MIPS 3 | 160 | 35 | 38 | $12 \%$ | $5.3 \times 2.1$ | $20 \times 2$ | $16 \times 18$ |  |



Fig. 6.7. MIPS filter transmissions, $R(\lambda)$. The dashed lines mark the reference wavelength for each band.

## Data Reduction and Photometry

The table of the available MIPS data for the DGS is given in Table 3.6 and in Bendo et al. (2012) who give a detailed description of the data reduction and of the photometry for integrated flux densities. Bendo et al. (2012) reprocessed the raw Spitzer data using the MIPS Data Analysis Tool (Gordon et al. 2005). The maps are given in units of $\mathrm{MJy} \mathrm{sr}^{-1}$. A non-linearity correction is applied to the $70 \mu \mathrm{~m}$ pixels that exceeded $66 \mathrm{MJy} \mathrm{sr}^{-1}$. This correction is given in Dale et al. (2007) based on data from Gordon et al. (2007):

$$
\begin{equation*}
F_{70}(\text { true })=0.581 \times F_{70}(\text { measured })^{1.13} \tag{6.17}
\end{equation*}
$$

The fluxes were extracted from the maps via aperture photometry. Bendo et al. (2012) used elliptical apertures with minor and major axes that were the greater of either 1.5 times the $\mathrm{D}_{25}$ isophotes of de Vaucouleurs et al. (1991) or 3'. This is to ensure that they do not need to apply aperture corrections on the $160 \mu \mathrm{~m}$ measurements. The whole group for Mrk1089 and UM311 was
taken into account in the aperture used for these sources as it is impossible to properly separate them at FIR wavelengths. Some of the DGS galaxies do not have an optical disk defined anywhere in the literature, so the $3^{\prime}$ aperture was used. For some very faint galaxies (fainter than 10 mJy at 24 $\mu \mathrm{m}$ ), $1^{\prime}$ apertures were used to minimize the effect of background noise in very large apertures (see Bendo et al. 2012). Special treatments were necessary for HS0052+2536, IC10, IIZw40, Mrk153, NGC5253, NGC6822, SBS1249+493, Tol0618-402 and Tol1214-277; and are detailed in Bendo et al. (2012).

Of the DGS sample, 43, 34 and 28 galaxies have been observed by MIPS at 24, 70 and $160 \mu \mathrm{~m}$ respectively (see Fig. 3.13). Bendo et al. (2012) MIPS flux densities compare well with previously published MIPS samples containing a subset of the DGS galaxies (Dale et al. 2007; Engelbracht et al. 2008). Therefore we are confident about the reliability of these results and will use them to perform the comparison with our PACS flux densities.

MIPS photometry for KINGFISH galaxies is presented in Dale et al. (2007).

## Comparison with PACS measurements for the DGS

The PACS flux densities correspond to monochromatic values for sources with spectra where $\nu F_{\nu}$ is constant, while the MIPS flux densities are monochromatic values for sources with the spectra of a $10^{4} \mathrm{~K}$ blackbody, so colour corrections need to be applied to measurements from both instruments before they are compared to each other. We first assume that the source spectral shape will be very close to that of the Rayleigh Jeans side of a modified blackbody: $\mathrm{F}_{S}(\nu) \propto \nu^{\alpha_{S}}$. We fit a modified blackbody through the three PACS data points to determine the spectral index $\alpha_{S}$, and apply the corresponding colour corrections from the available PACS documentation ${ }^{5}$. For the MIPS flux densities, we fit a blackbody through the 70 and $160 \mu \mathrm{~m}$ data points (not using the $24 \mu \mathrm{~m}$ point) and apply the corrections from the MIPS Handbook ${ }^{6}$. The typical colour corrections for MIPS are of the order of $10 \%$ and $4 \%$ on average at 70 and $160 \mu \mathrm{~m}$. However, they are of the order of $1 \%$ or $2 \%$ in the 70 and $160 \mu \mathrm{~m}$ PACS bands. For non detected galaxies, where we, for PACS, and/or Bendo et al. (2012), for MIPS, reported upper limits (9 galaxies), we are not able to properly fit a blackbody and therefore derive a proper colour correction. We do not compare PACS and MIPS flux densities for these galaxies for now.

We use the ratios of the PACS and MIPS flux densities to assess how well the measurements from the instrument agree with each other; a ratio of 1 corresponds to a very good agreement. The PACS/MIPS ratios at 70 and $160 \mu \mathrm{~m}$ are shown in Figure 6.8, and the correspondence is relatively good. The PACS/MIPS ratio is $1.019 \pm 0.112$ at $70 \mu \mathrm{~m}$ and $0.995 \pm 0.153$ at $160 \mu \mathrm{~m}$. This is to be compared to an average uncertainty of $\sim 12 \%(\sim 11 \%$ from MIPS and $\sim 5 \%$ for PACS, added in quadrature) and $\sim 16 \%$ ( $\sim 15 \%$ from MIPS and $\sim 7 \%$ for PACS, added in quadrature) on the ratios at 70 and $160 \mu \mathrm{~m}$ respectively.

If we now consider galaxies detected at $70 \mu \mathrm{~m}$ and not at $160 \mu \mathrm{~m}$, indicated by a different symbol on the left panel of Figure 6.8, we are still able to compare, with extra caution, the measurements at $70 \mu \mathrm{~m}$. Indeed, as we are not able to derive a proper colour correction for those galaxies, we add to the MIPS $70 \mu$ m flux densities a $10 \%$ uncertainty and a $1 \%$ uncertainty to the PACS $70 \mu \mathrm{~m}$ flux densities to account for the colour correction effect. When adding these extra galaxies at $70 \mu \mathrm{~m}$, the PACS/MIPS ratio is $0.985 \pm 0.158$ at $70 \mu \mathrm{~m}$. This is to compare with an average uncertainty of $\sim 14 \%$ on the $70 \mu \mathrm{~m}$ ratio ( $\sim 12 \%$ from MIPS and $\sim 7 \%$ for PACS, added in quadrature, including the extra galaxies). The very faint and discrepant galaxies at $70 \mu \mathrm{~m}$ are HS1222+3741 (ratio of 0.40)

[^29]

Fig. 6.8. Comparison of PACS flux densities and MIPS flux densities: PACS-to-MIPS flux density ratios as a function of PACS flux density at $70 \mu \mathrm{~m}$ (left) and $160 \mu \mathrm{~m}$ (right). As a guide to the eye, the unity line is added as a solid line as well as the average uncertainties on the ratio in both bands as dotted lines. These average uncertainties are $\sim 12 \%$ and $\sim 16 \%$ at 70 and $160 \mu \mathrm{~m}$. Colours distinguish the selected mapping method (see Chapter 4).
and Tol1214-277 (ratio of 0.24 ). For HS1222+3741, the MIPS image contains some bright pixels near the edge of the photometry aperture used for MIPS, and this may have driven the $70 \mu \mathrm{~m}$ MIPS flux density up. For Tol1214-277, a nearby source is present in the MIPS data and, although its contribution has been subtracted when computing the MIPS $70 \mu \mathrm{~m}$ flux, some contribution from this source may still be present. Additionally, measuring accurate flux densities at $\leq 50 \mathrm{mJy}$ in both MIPS and PACS data is difficult and may have led to the discrepancies.

The error on the average ratio of the PACS to MIPS flux densities is comparable to the average uncertainties on the ratio for both bands. Thus there is a good photometric agreement between PACS and MIPS photometry for the DGS sample.

The same test has been done by Dale et al. (2012) for the KINGFISH galaxies. They found that the Spitzer and Herschel flux densities were also in good agreement.

### 6.3 WISE

## Instrument

The Wide-field Infrared Survey Explorer (WISE, Wright et al. 2010) surveyed the entire sky at 3.4, 4.6, 12 and $22 \mu \mathrm{~m}$. The telescope contains four mid-infrared focal plane array detectors that observed the sky simultaneously in a $47^{\prime} \times 47^{\prime} \mathrm{FOV}$. The main characteristics for WISE are given in Table 6.5 and the 4 broad-band MIR filters are shown in Fig. 6.9. The spectral convention for WISE is $\mathrm{F}_{S}(\nu) \propto \nu^{-2}$ and thus Eq. 6.9 gives:

$$
\begin{equation*}
F_{S}^{\prime}\left(\nu_{0}\right)=\frac{\int F_{S}^{\prime}(\nu)\left(\nu_{0} / \nu\right) R(\nu) \mathrm{d} \nu}{\int\left(\nu_{0} / \nu\right)^{3} R(\nu) \mathrm{d} \nu} \tag{6.18}
\end{equation*}
$$

Table 6.5. WISE characteristics.

| WISE |  |  |  |  |
| :--- | :---: | :---: | :---: | :---: |
| Band | Wavelength $(\mu \mathrm{m})$ | Bandwidth $(\mu \mathrm{m})$ | FWHM $\left({ }^{\prime \prime}\right)$ | $F_{0}[\mathrm{Jy}]$ |
| WISE 1 | 3.4 | 0.750 | 6.1 | 306.682 |
| WISE 2 | 4.6 | 1.015 | 6.4 | 170.663 |
| WISE 3 | 12.0 | 1.425 | 6.5 | 29.045 |
| WISE 4 | 22.0 | 2.905 | 12.0 | 8.284 |

## Photometry: Searching the literature

The WISE data have been extracted from the NASA/IPAC ISA database and the WISE all-sky catalog ${ }^{7}$. Magnitudes are given, and can be converted into flux densities using Eq. 6.1 and the zero-magnitude flux values from Wright et al. (2010), reported in Table 6.5.

According to the WISE documentation ${ }^{8}$, the optimal flux measurement for point sources is the $\mathrm{w}_{i}$ MPRO (with $\mathrm{i}=\{1,2,3,4\}$ ) measurement from the database, as it comes from profile-fitting photometry. However when the source is resolved or extended, it is better to use the circular aperture photometry reported: $\mathrm{W}_{i} \mathrm{MAG}$. This is indicated by an EXTFLG value of 5 in the database. The WISE flux densities for 40 DGS galaxies are reported in Table 6.10.


Fig. 6.9. WISE filter transmissions, $R(\lambda)$. The dashed lines mark the reference wavelength for each band.

### 6.4 IRAS

The InfraRed Astronomical Satellite (IRAS) was designed to observe the entire sky in four wavelengths bands centred on $12,25,60$ and $100 \mu \mathrm{~m}$. The basic characteristics of $I R A S$ are given in Table 6.6 and the four broadband filters are shown in Fig. 6.10. IRAS data for the DGS have been compiled from the literature: the NASA/IPAC ISA IRAS Faint Source (v2.0 1990) and Point

[^30]Source (v2.1) Catalogs, Rice et al. (1988); Sanders et al. (2003); Engelbracht et al. (2008) and is given in Table 6.11. The spectral convention for $\operatorname{IR} A S$ is $\mathrm{F}_{S}(\nu) \propto \nu^{-1}$ and thus Eq. 6.9 gives:

$$
\begin{equation*}
F_{S}^{\prime}\left(\nu_{0}\right)=\frac{\int F_{S}^{\prime}(\nu)\left(\nu_{0} / \nu\right) R(\nu) \mathrm{d} \nu}{\int\left(\nu_{0} / \nu\right)^{2} R(\nu) \mathrm{d} \nu} \tag{6.19}
\end{equation*}
$$

Table 6.6. $I R A S$ characteristics.

| IRAS |  |  |  |
| :--- | :---: | :---: | :---: |
| Band | $\lambda(\mu \mathrm{m})$ | Bandwidth $(\mu \mathrm{m})$ | Calibration accuracy |
| IRAS 1 | 12 | 4.7 | $2 \%$ |
| IRAS 2 | 25 | 19 | $5 \%$ |
| IRAS 3 | 100 | 35 | $5 \%$ |
| IRAS 4 | 160 | 35 | $10 \%$ |



Fig. 6.10. IRAS filter transmissions, $R(\lambda)$. The dashed lines mark the reference wavelength for each band.

Note that $\operatorname{IRAS}$ is not very sensitive and that the resolution varies from $\sim 30^{\prime \prime}$ at $12 \mu \mathrm{~m}$ to $\sim 120^{\prime \prime}$ at $100 \mu \mathrm{~m}$. Even for the detected sources, such a low resolution may imply that several sources are mixed in the beam and indistinguishable. Thus some of the $I R A S$ fluxes of Table 6.11 may be unreliable and $I R A S$ photometry should be treated with caution. 25 galaxies have IRAS measurements and only 15 are detected in all of the IRAS bands.

### 6.5 Ground-based submm data

### 6.5.1 Instruments

SCUBA The Submillimeter Common-User Bolometer Array (SCUBA) is a submm bolometre receiver, mounted on the JCMT (James Clerk Maxwell Telescope) in Hawaii, composed of two arrays: the Long-Wave (LW) array observing the sky at $850 \mu \mathrm{~m}$ and the Short-Wave (SW) array observing the sky at $450 \mu \mathrm{~m}$. The FOV is $2.3^{\prime}$ in diameter and the sky is observed simultaneously
by the two arrays, with FWHM of $7.5^{\prime \prime}$ and $14^{\prime \prime}$ for the SW and LW arrays respectively. SCUBA was upgraded to SCUBA2 in the Fall 2011.

LABOCA The Large APEX BOlometer CAmera (LABOCA) is one of the instruments mounted on APEX (Atacama Pathfinder EXperiment) in Chile, and was designed by the MPIfR. It is a multi-beam bolometric receiver observing in the submillimetric atmospheric window at $870 \mu \mathrm{~m}$ with a FOV of $11.4^{\prime} \times 11.4^{\prime}$ and a beam FWHM of $19.5^{\prime \prime}$.

MAMBO MAMBO (the MAx-planck Millimeter-BOlometer) was a submm/mm bolometre array, mounted on the IRAM $30-\mathrm{m}$ telescope in Spain, observing at 1.2 mm with a FOV of $4^{\prime} \times 4^{\prime}$ and a FWHM of 11". MAMBO has been decommissioned in April 2011.

The main characteristics for these three instruments are presented in Table 6.7 and the bandpass filters are shown in Fig. 6.11.

Table 6.7. SCUBA, LABOCA and MAMBO characteristics.

| Band | $\lambda[\mu \mathrm{m}]$ | Bandwidth $[\mu \mathrm{m}]$ | FWHM $\left({ }^{(\prime \prime}\right)$ | FOV $\left({ }^{\prime} \times{ }^{\prime}\right)$ |
| :--- | :---: | :---: | :---: | :---: |
| SCUBA SW $^{a}$ | 450 | $90^{d}$ | 7.5 | 2.3 (diameter) |
| SCUBA LW $^{a}$ | 850 | $170^{d}$ | 14 | 2.3 (diameter) |
| LABOCA $^{b}$ | 870 | 160 | $19.5^{e}$ | $11.4 \times 11.4$ |
| MAMBO $^{c}$ | 1200 | 430 | 11 | $4 \times 4$ |

${ }^{a}$ : Values from: http://www.jach.hawaii.edu/JCMT/continuum/scuba/intro.html.
${ }^{b}$ : Values from: http://www3.mpifr-bonn.mpg.de/staff/gsiringo/laboca/laboca_at_the_mpifr_ bolometer_group.html.
${ }^{c}$ : Values from: http://www.iram.es/IRAMES/mainWiki/PoolObserving.
${ }^{d}$ : Holland et al. (1999).
${ }^{e}$ : Albrecht et al., in prep.


Fig. 6.11. Ground-based submm filter transmissions, $R(\lambda)$. The dashed lines mark the reference wavelength for each band.

### 6.5.2 Literature measurements

SCUBA data for 21 KINGFISH galaxies ( $34 \%$ of the sample) is presented in Dale et al. (2007). We find submm literature measurements for 11 galaxies of the DGS: Haro11, Mrk1089, NGC1705 and UM311 with LABOCA (Galametz et al. 2009), Haro2 with SCUBA Dale et al. (2007), He2-10, IIZw40, NGC1140, NGC1569, NGC4214 and NGC4449 with SCUBA and MAMBO (Lisenfeld et al. 2002; Böttner et al. 2003; Galliano et al. 2003, 2005; Kiuchi et al. 2004; Hermelo et al. 2013).

We are interested here in the analysis of the global dust properties of galaxies. For NGC4214 and NGC4449 we do not use SCUBA data as the field of view of the observations is very small and the cold diffuse dust emission is missed by the observations: the maps presented in Kiuchi et al. (2004) for NGC4214 only show the centre of the galaxy in a $2^{\prime} \times 1.40^{\prime}$ field of view, and Böttner et al. (2003) present maps of NGC4449 on a $2.8^{\prime} \times 3^{\prime}$ field of view. Galametz et al. (2009) report a flux density at $870 \mu \mathrm{~m}$ for UM311 corresponding to the regions 12 and 3 of Moles et al. (1994) and not for the full group that we are considering. Thus we do not use this flux density either.

### 6.5.3 New LABOCA observations

In order to complete this sample of submm observations, we wrote 8 proposals for the three instruments: one for MAMBO, two for SCUBA2 and five for LABOCA, out of which 7 were accepted. However MAMBO was decommissioned right after the acceptance of our proposal and thus no observations could be carried on. For SCUBA2 observations, because of bad weather conditions and a medium priority for our program, we were not able to observe the proposed sources down to the required rms. For LABOCA the strategy was to separate our requested time between two of the owners of the APEX telescope time: ESO (European Southern Observatory) and OSO (Onsala Space Observatory, Sweden). Our proposals were accepted and we were able to successfully observe three sources out of the four requested (HS0052+2536 was dropped because it was not detected after all of the allocated time). Thus we present here LABOCA observations for Mrk930, NGC5253 and UM448.

## Observations and data reduction

The sources were observed with LABOCA on APEX between May 2012 and May 2013 from programs 089.B0922, 089.F9306, 091.B0724 and 091.F9306. The sources were mapped using the spiral raster map mode with map sizes adjusted from the extent of the SPIRE $250 \mu \mathrm{~m}$ emission as the two beam sizes are comparable ( $\sim 18-19^{\prime \prime}$ ). The focus settings of the telescope were regularly checked during the observations using planets (Uranus and Neptune). The pointing of the telescope was checked on strong sources located close to the target galaxy. Skydips were performed regularly during the observations in order to derive the zenith opacity needed to correct the observations for atmospheric attenuation. Calibration of the flux densities was derived mostly from observations of planets (Mars, Uranus and Neptune as primary calibrators). The calibration was found to be accurate to $12 \%$.

The data reduction is performed with the BoA software package ${ }^{9}$ (BOlometer Array analysis software, Schuller 2012). The following steps are applied to the raw time series:

- calibration correction;
- opacity correction using the interpolation of the opacity at the elevation of the scan (from the skydips);
- correction for the thermal drift of the bolometres;

[^31]- flat fielding;
- conversion from counts to Jy;
- flagging of bad channels: Additionally to the known dead and noisy channels, all channels that deviate from the median rms of all of the channels were flagged after checking all of the time series to identify them. After this step, about $83 \%$ of the total 295 channels remains unflagged;
- removal of correlated noise on the full array and on groups of channels sharing the same amplifier box and the same wiring (this noise on groups of channels was found during the commissioning period of LABOCA);
- baseline subtraction;
- despiking;
- low-frequency filtering;
- flagging of the data taken outside reasonable telescope scanning velocity and acceleration limits.

Then each reduced scan is gridded in a rms-weighted intensity map with $4^{\prime \prime} \times 4^{\prime \prime}$ pixels. Individuals maps are coadded and noise-weighted again to derive the final maps. As our sources are extended, it is important to differentiate between correlated noise and extended emission. However the "blind" skynoise removal step in the data reduction can not separate correlated atmospheric noise from the real diffuse extended emission of the sources. To avoid this effect, an iterative method is applied where a model of the flux distribution of the source is determined from the observations and improved at each new iteration of the data reduction. This data reduction method and the reduction parameters have been tested and improved in order to get the best recovery of the faint diffuse emission of extended sources. The data reduction is presented in more details in Albrecht et al., (in prep.) on another sample of galaxies. The new LABOCA maps are presented in Fig. 6.12. The rms level in the maps is $0.04 \mathrm{mJy} / \mathrm{beam}$ for Mrk930 and NGC5253 and $0.05 \mathrm{mJy} /$ beam for UM448.

## Photometry

Aperture photometry is performed on the maps to derive flux densities. Maps are first converted from mJy/beam to mJy/pix using an effective beam area of $521 \operatorname{arcsec}^{2}$ (Albrecht et al. in prep.). For NGC5253, we use the same aperture as for Herschel measurements. For Mrk930 and UM448, the aperture used for Herschel photometry encompass a small fraction of nearby sources (to the east of Mrk930 and to the north of UM448, see Fig. 6.12). For these two sources, we adapt the shape of the Herschel aperture to only encompass the emission of the galaxy at $870 \mu \mathrm{~m}$. The Herschel background aperture is used for NGC5253. For Mrk930 and UM448, the nearby sources may bias our estimation of the background if we take an annulus around the source aperture. Thus we use small circular apertures of $30^{\prime \prime}$ radius spread around the galaxy for these two sources. Errors on the flux density are estimated by adapting the method presented in Chap 4 to LABOCA.

We get a total flux density at $870 \mu \mathrm{~m}$ for the sources of $F_{870}=41 \pm 5 \mathrm{mJy}$ for Mrk930, $F_{870}=494 \pm 59 \mathrm{mJy}$ for NGC5253 and $F_{870}=76 \pm 9 \mathrm{mJy}$ for UM448. These flux densities are reported in Table 6.12. If we use the apertures used previously for Herschel photometry, we get $F_{870}=42 \pm 6 \mathrm{mJy}$ for Mrk 930 and $F_{870}=75 \pm 9 \mathrm{mJy}$ for UM448 which is consistent with our results.

James et al. (2002) observed NGC5253 and UM448 at $850 \mu \mathrm{~m}$ with SCUBA and reported flux densities at $850 \mu \mathrm{~m}$ of $192 \pm 23 \mathrm{mJy}$ for NGC5253 ( $39 \%$ of our value) and $40 \pm 9 \mathrm{mJy}$ for UM448 ( $53 \%$ of our value). We believe that these discrepancies are due to the apertures used by James et al. (2002) which are much smaller than ours: $41^{\prime \prime}$ in diameter for NGC5253 and $34^{\prime \prime}$ in diameter for UM448. Indeed when using the apertures from James et al. (2002), we find a flux density at 870 $\mu \mathrm{m}$ of $212 \pm 25 \mathrm{mJy}$ for NGC5253 and $29.6 \pm 3.4 \mathrm{mJy}$ for UM448, much closer to their findings at $850 \mu \mathrm{~m}$. Additionally, our data reduction paid particular attention to recover the diffuse extended emission whereas James et al. (2002) do not mention any particular treatment of the extended emission in their galaxies.


Fig. 6.12. LABOCA $870 \mu \mathrm{~m}$ maps for Mrk930 (top left), NGC5253 (top right) and UM448 (bottom), in units of $\mathrm{MJy} / \mathrm{sr}$. The contours are $\mathrm{S} / \mathrm{N}$ contours with $\mathrm{S} / \mathrm{N}=3,5$ and 6 for Mrk930, $\mathrm{S} / \mathrm{N}=3,5,6,10,20$, $30,50,80$ for NGC5253 and $\mathrm{S} / \mathrm{N}=3,5,6,10$ for UM448. The black cross marks the position of the galaxy from NED and the LABOCA beam is shown on the bottom right corner of the maps.

### 6.5.4 Non-dust contaminations

Possible contamination from non-dust sources can occur at submm wavelengths and we need to take them into account when using the submm flux densities to constrain a SED dust model. We can identify two sources of contamination at our wavelengths of interest:

Radio emission: There are two sources of radio emission: thermal (bremsstrahlung, also called free-free) and non-thermal (synchrotron) emission. Free-free emission comes from charged particles when they are accelerated during collisions. It arises from the interaction of very fast charged particles with a gas and is emitted from the ionised phase of the ISM (e.g., Hir regions, see Section 1.3). This emission can be estimated using the radio tendency: $L_{\nu}^{f f} \propto \nu^{-0.1}$ or emission from the $\mathrm{H} \alpha$ recombination line. Synchrotron emission arises when relativistic electrons are gyrating about a magnetic field. The synchrotron emission is also well described by a power law of spectral index $\alpha_{s y n c}: L_{\nu}^{s y n c} \propto \nu^{-\alpha_{s y n c}}$. A typical value for $\alpha_{s y n c}$ is $0.8 \pm 0.1$. Usually, if no strong AGN is present, the free-free component dominates over the synchrotron component of the radio emission. For a review on radio emission in normal galaxies, see the review by Condon (1992).

CO emission: Some CO transitions can fall within the wavelength range we cover: $\mathrm{CO}(2-1)$ at 1.3 mm in the MAMBO 1.2 mm band, $\mathrm{CO}(3-2)$ at $867 \mu \mathrm{~m}$ in the LABOCA $870 \mu \mathrm{~m}$ and the SCUBA $850 \mu \mathrm{~m}$ bands. $\mathrm{CO}(1-0)$ can be difficult to detect in low-metallicity dwarf galaxies (e.g. Leroy et al. 2005; Schruba et al. 2012), but, in principle, other higher excitation states could be present (Cormier et al. 2014).

Estimating the contaminations: To estimate the contamination from radio emission, we assume that we can neglect synchrotron emission compared to the free-free emission. We collect at least one radio flux density for each galaxy and extrapolate to 850,870 and $1200 \mu \mathrm{~m}$ using the power-law description of the radio emission: $L_{\nu}^{\text {radio }} \propto \nu^{-0.1}$. For some galaxies, several radio flux densities are available. In this case the radio contamination from each available flux density is estimated and we use the mean and standard deviation of all of the estimated contaminations for the final radio contamination. The mean radio contamination to the total flux density at $850 \mu \mathrm{~m}$ for the DGS galaxies is $15.8 \%$ and $16.8 \%$ for the KINGFISH sample (giving $16.6 \%$ for the total sample). At $870 \mu \mathrm{~m}$ and $1200 \mu \mathrm{~m}$ respectively the mean radio contaminations are $15.3 \%$ and $22.5 \%$.

The CO emission in the submm filters can be estimated using:

$$
\begin{equation*}
F_{C O}=\frac{2 k \nu^{3}}{c^{3} \Delta \nu} \int_{\Omega_{s}} I_{C O} \mathrm{~d} \Omega \tag{6.20}
\end{equation*}
$$

in the case where a CO map is available (for the most extended sources). $k$ is the Boltzmann constant, $\nu$ is the frequency of the CO transition, $\Delta \nu$ is the bandwidth of the submm bolometre ( 70 GHz for SCUBA at $850 \mu \mathrm{~m}, 60 \mathrm{GHz}$ for LABOCA and 80 GHz for MAMBO), $\Omega_{s}$ is the solid angle covering the source and $I_{C O}$ is the velocity integrated intensity given by: $I_{C O}=\int T_{m b} \mathrm{~d} v$ where $T_{m b}$ is the main beam temperature. In some cases, CO luminosities, $L_{C O}$, are given using the distance $D$ to the source. We convert these luminosities using: $\int_{\Omega_{s}} I_{C O} \mathrm{~d} \Omega=L_{C O} / D^{2}$.

For compact sources, where a single pointing encompasses the whole galaxy, Eq. 6.20 transforms into:

$$
\begin{equation*}
F_{C O}=\frac{2 k \nu^{3}}{c^{3} \Delta \nu} I_{C O} \Omega_{m b} \tag{6.21}
\end{equation*}
$$

where the solid angle $\Omega_{m b}$ is given by $\frac{\pi \theta^{2}}{4 \ln (2)}$ with $\theta$ the FWHM of the main beam. In some cases, $\mathrm{CO}(3-2)$ or $\mathrm{CO}(2-1)$ measurements are not available to directly estimate the CO contribution to the

850 and $870 \mu \mathrm{~m}$, and $1200 \mu \mathrm{~m}$ bands respectively. For these galaxies, we use $\mathrm{CO}(2-1)$ or $\mathrm{CO}(1-0)$ in lieu of $\mathrm{CO}(3-2)$, and $\mathrm{CO}(1-0)$ in lieu of $\mathrm{CO}(2-1)$ if lower J CO transition measurements are available. We convert the CO intensities of the lower J transitions to the higher J transitions using the band ratios from Wilson et al. (2012) obtained on a sample of 47 galaxies:

$$
\left\{\begin{array}{l}
R_{3-2 / 2-1}=0.36 \pm 0.04  \tag{6.22}\\
R_{3-2 / 1-0}=0.18 \pm 0.02 \\
R_{2-1 / 1-0}=0.50 \pm 0.08
\end{array}\right.
$$

These lines ratios are relatively consistent too with the line ratios derived by Cormier et al. (2014) on a sample of 5 dwarf galaxies. We did not find any CO measurement in the literature for NGC1705 and we use the mean CO line contamination to the submm flux densities. This mean CO line contamination to the total flux density at $850 \mu \mathrm{~m}$ for the DGS galaxies is $2.9 \%$ and $7.8 \%$ for the KINGFISH sample (giving $6.8 \%$ for the total sample). At $870 \mu \mathrm{~m}$ and $1200 \mu \mathrm{~m}$ respectively the mean line contaminations are $0.8 \%$ and $1.7 \%$. When we have CO non detections, we consider the upper limit for the CO line contamination to be conservative.

The final submm flux densities, presented in Tables 6.12 and 6.13, are derived by removing both the radio and CO line contaminations from the original flux densities. The mean radio+CO contamination to the total flux density at $850 \mu \mathrm{~m}$ for the DGS galaxies is $18.7 \%$ and $24.5 \%$ for the KINGFISH sample (giving $23.4 \%$ for the total sample). At $870 \mu \mathrm{~m}$ and $1200 \mu \mathrm{~m}$ respectively the mean radio + CO contaminations are $16.1 \%$ and $24.3 \%$. Note that the lower mean contaminations in the $870 \mu \mathrm{~m}$ for the radio, CO , or radio +CO contaminations, compared to the same contaminations at $850 \mu \mathrm{~m}$, is only representative of the smaller number of sources observed at $870 \mu \mathrm{~m}$ ( 6 galaxies) compared to the number of sources observed at $850 \mu \mathrm{~m}$ (26 galaxies).

Table 6.8. 2MASS flux densities for the DGS galaxies. When an upper limit is given, it is the $95 \%$ confidence upper limit given by the database.

| Source | $\mathrm{F}_{J}$ [mJy] | $\sigma_{F_{J}}[\mathrm{mJy}]$ | $\mathrm{F}_{H}$ [mJy] | $\sigma_{F_{H}}[\mathrm{mJy}]$ | $\mathrm{F}_{K_{s}}[\mathrm{mJy}]$ | $\sigma_{F_{K_{s}}}[\mathrm{mJy}]$ | Ref |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Haro11 | 13.0 | 0.4 | 13.0 | 0.6 | 14.1 | 0.7 | 1 |
| Haro2 | 43.6 | 1.4 | 53.5 | 2.3 | 44.9 | 2.3 | 2 |
| Haro3 | 44.7 | 1.4 | 53.2 | 2.3 | 37.9 | 1.9 | 1 |
| He2-10 | 178 | 6 | 201 | 9 | 167 | 8 | 1 |
| HS0017+1055 | 0.38 | 0.01 | 0.59 | 0.03 | 0.36 | 0.02 | 3 |
| HS0052+2536 | 0.54 | 0.02 | 0.55 | 0.02 | 0.40 | 0.02 | 3 |
| HS0822+3542 | 0.15 | 0.02 | 0.34 | 0.04 | 0.16 | 0.03 | 4 |
| HS1222+3741 | 0.90 | 0.03 | 1.18 | 0.05 | 0.83 | 0.04 | 3 |
| HS1236+3937 | - | - | - | - | - | - | 3 |
| HS1304+3529 | 0.32 | 0.01 | 0.27 | 0.01 | $\leq 0.385$ | - | 3 |
| HS1319+3224 | - | - | - | - | - | - | 3 |
| HS1330+3651 | - | - | - | - | - | - | 3 |
| HS1442+4250 | - | - | - | - | - | - | 3 |
| HS2352+2733 | - | - | - | - | - | - | 3 |
| IZw18 | 0.75 | 0.08 | 0.64 | 0.07 | 0.52 | 0.06 | 3 |
| IC10 | 2052 | 64 | 2972 | 129 | 2632 | 133 | 2 |
| IIZw40 | 15.1 | 1.5 | 21.0 | 2.2 | 20.3 | 2.1 | 4 |
| Mrk1089 | 12.2 | 0.4 | 13.9 | 0.6 | 12.2 | 0.6 | 1 |
| Mrk1450 | 1.05 | 0.03 | 0.90 | 0.04 | 0.66 | 0.03 | 3 |
| Mrk153 | 4.64 | 0.56 | 4.27 | 0.59 | 3.63 | 0.51 | 4 |
| Mrk209 | 7.42 | 1.34 | 8.22 | 2.00 | 7.28 | 2.40 | 5 |
| Mrk930 | 3.52 | 0.36 | 4.64 | 0.48 | 3.53 | 0.36 | 4 |
| NGC1140 | 48.9 | 1.5 | 57.2 | 2.5 | 41.8 | 2.1 | 2 |
| NGC1569 | 473 | 49 | 547 | 57 | 479 | 50 | 2 |
| NGC1705 | 52.1 | 1.6 | 50.9 | 2.2 | 41.1 | 2.1 | 2 |
| NGC2366 | 145 | 10 | 147 | 13 | 110. | 14. | 5 |
| NGC4214 | 521 | 16 | 614 | 27 | 458 | 23 | 2 |
| NGC4449 | 916 | 29 | 1068 | 46 | 839 | 42 | 2 |
| NGC4861 ${ }^{a}$ | 16.9 | 0.5 | 14.0 | 0.6 | 13.1 | 0.7 | 2 |
| NGC5253 | 380 | 12 | 414 | 18 | 334 | 17 | 2 |
| NGC625 | 218 | 7 | 236 | 10 | 184 | 9 | 2 |
| NGC6822 ${ }^{a}$ | 1619 | 50 | 1754 | 76 | 1363 | 69 | 2 |
| Pox186 | - | - | - | - | - | - | 3 |
| SBS0335-052 | 0.30 | 0.01 | 0.28 | 0.01 | $\leq 0.38$ | - | 3 |
| SBS1159+545 | - | - | - | - | - | - | 3 |
| SBS1211+540 | - | - | - | - | - | - | 3 |
| SBS1249+493 | - | - | - | - | - | - | 3 |
| SBS1415+437 | - | - | - | - | - | - | 3 |
| SBS1533+574 | 1.95 | 0.06 | 2.14 | 0.09 | 1.54 | 0.08 | 3 |
| Tol0618-402 | 4.32 | 0.13 | 5.84 | 0.25 | 4.04 | 0.20 | 4 |
| Tol1214-277 | - | - | - | - | - | - | 3 |
| UGC4483 | 2.20 | 0.23 | 0.94 | 0.12 | 3.11 | 0.39 | 4 |
| UGCA20 | - | - | - | - | - | - | 3 |
| UM133 | - | - | - | - | - | - | 3 |
| UM311 | - | - | - | - | - | - | 3 |
| UM448 | 19.3 | 0.6 | 22.2 | 1.0 | 19.4 | 1.0 | 2 |
| UM461 | 2.03 | 0.21 | 1.49 | 0.16 | 1.27 | 0.14 | 4 |
| VIIZw403 | 4.31 | 0.44 | 3.89 | 0.40 | 2.97 | 0.31 | 4 |

${ }^{a}$ : The aperture used does not cover the total emission from the galaxy and the magnitudes reported in the database are not consistent with the rest of the IR photometry. Thus we do not consider them for the modelling.
References: (1) 2MASS Extended Objects Final Release 2003; (2) Jarrett et al. (2003) ; (3) NASA/IPAC ISA Point Source Catalog, available at http://irsa.ipac.caltech.edu/cgi-bin/Gator/nph-dd?catalog=fp_psc ; (4) Engelbracht et al. (2008) ; (5) Dale et al. (2009)
Table 6.9. IRAC flux densities for the DGS galaxies. When an upper limit is given, it is the $3 \sigma$ upper limit computed in Section 6.2 .1 Photometry.

| Source | $\mathrm{F}_{3.6}$ [mJy] | $\sigma_{3.6}[\mathrm{mJy}]$ | $\mathrm{F}_{4.5}$ [mJy] | $\sigma_{4.5}[\mathrm{mJy}]$ | $\mathrm{F}_{5.8}$ [mJy] | $\sigma_{5.8}[\mathrm{mJy}]$ | $\mathrm{F}_{8.0}$ [mJy] | $\sigma_{8.0}[\mathrm{mJy}]$ | Radius (") | Aper corr ? |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Haro11 | 22.4 | 3.2 | 32.3 | 4.6 | 77.1 | 10.9 | 164 | 23 | $45^{a}$ | yes |
| Haro2 | 24.7 | 3.5 | 18.2 | 2.6 | 49.3 | 7.0 | 257 | 36 | $50^{a}$ | yes |
| Haro3 | 28.2 | 4.0 | 21.5 | 3.1 | 43.9 | 6.3 | 116 | 16 | $60^{a}$ | yes |
| He2-10 | 134 | 19 | 99.2 | 14.1 | 296 | 42 | 836 | 118 | $108^{a}$ | yes |
| HS0017+1055 | 0.17 | 0.02 | 0.31 | 0.03 | 0.76 | 0.08 | 1.53 | 0.16 | $12^{\text {b }}$ |  |
| HS0052+2536 | 0.72 | 0.12 | 0.32 | 0.09 | 0.66 | 0.17 | 3.17 | 0.46 | $17^{\text {c }}$ | yes |
| HS0822+3542 | 0.12 | 0.01 | 0.10 | 0.01 | 0.10 | 0.03 | 0.09 | 0.03 | $5^{b}$ |  |
| HS1222+3741 | 0.25 | 0.03 | 0.19 | 0.02 | 0.27 | 0.05 | 0.57 | 0.07 | $14^{a}$ |  |
| HS1236+3937 | $\leq 0.56$ | - | $\leq 0.38$ | - | $\leq 0.66$ | - | $\leq 0.32$ | - | $15^{a}$ |  |
| HS1304+3529 | 0.38 | 0.08 | 0.41 | 0.09 | $\leq 0.78$ | - | 0.60 | 0.14 | $18^{a}$ |  |
| HS1319+3224 | 0.06 | 0.01 | 0.05 | 0.01 | 0.10 | 0.02 | 0.12 | 0.02 | $8^{a}$ |  |
| HS1330+3651 | 0.56 | 0.09 | 0.46 | 0.10 | $\leq 0.80$ | - | 0.97 | 0.17 | $20^{a}$ |  |
| HS1442+4250 | - | - | 1.24 | 0.21 | - | - | 1.59 | 0.25 | $51^{a}$ | yes |
| HS2352+2733 | $\leq 0.04$ | - | $\leq 0.04$ | - | $\leq 0.13$ | - | $\leq 0.11$ | - | $15^{a}$ |  |
| IZw18 | 0.36 | 0.05 | 0.35 | 0.05 | 0.39 | 0.07 | 0.53 | 0.08 | $12^{\text {b }}$ | yes |
| IC10 ${ }^{f}$ | - | - | - | - | - | - | - | - | - | - |
| IIZw40 | 18.1 | 4.3 | 19.9 | 3.5 | 39.8 | 5.7 | 105 | 15 | $45 \times 33(39)^{c}$ | yes |
| Mrk1089 | 20.4 | 2.9 | 15.5 | 2.2 | 31.4 | 4.4 | 82.2 | 11.6 | $75^{a}$ | yes |
| Mrk1450 | 1.22 | 0.17 | 1.01 | 0.14 | 0.93 | 0.14 | 2.06 | 0.29 | $20^{a}$ | yes |
| Mrk153 | 2.29 | 0.33 | 1.70 | 0.24 | 1.76 | 0.25 | 2.75 | 0.39 | $25^{\text {b }}$ | yes |
| Mrk209 | 2.90 | 0.41 | 2.21 | 0.32 | 1.78 | 0.27 | 2.20 | 0.32 | $39^{\text {d }}$ | yes |
| Mrk930 | 2.66 | 0.38 | 2.73 | 0.39 | 4.17 | 0.59 | 9.65 | 1.37 | $13^{b}$ | yes |
| NGC1140 | 32.8 | 4.8 | 23.3 | 3.4 | 41.0 | 5.8 | 93.9 | 13.3 | $118^{\text {a }}$ | yes |
| NGC1569 | 312 | 45 | 244 | 35 | 324 | 46 | 533 | 75 | $150^{a}$ | yes |
| NGC1705 | 25.5 | 5.9 | 18.3 | 4.2 | 15.0 | 2.8 | 17.5 | 2.6 | $72^{a}$ | yes |
| NGC2366 | 61.5 | 12.4 | 45.7 | 10.4 | 35.8 | 7.6 | 52.8 | 8.1 | $200^{\text {d }}$ | yes |
| NGC4214 | 334 | 47 | 253 | 36 | 280 | 40 | 748 | 106 | $205 \times 160(181)^{c}$ | yes |
| NGC4449 | 472 | 67 | 342 | 48 | 709 | 100 | 1622 | 229 | $190 \times 170(180)^{c}$ | yes |
| NGC4861 | 18.3 | 2.6 | 13.8 | 2.0 | 16.8 | 2.5 | 17.7 | 2.5 | $150 \times 100(71)^{c}$ | yes |
| NGC5253 | 235 | 34 | 252 | 36 | 458 | 65 | 812 | 115 | $120^{a}$ | yes |
| NGC625 | 110 | 16 | 79.7 | 11.4 | 71.3 | 10.1 | 147 | 21 | $170^{a}$ | yes |
| NGC6822 | 1780 | 272 | 1240 | 196 | 951 | 141 | 1226 | 175 | $440{ }^{\text {a }}$ | yes |
| Pox186 | 0.18 | 0.02 | 0.18 | 0.02 | 0.14 | 0.03 | 0.64 | 0.07 | $9^{b}$ |  |
| SBS0335-052 | 0.78 | 0.08 | 1.64 | 0.17 | 4.50 | 0.45 | 12.3 | 1.2 | $10^{a}$ |  |
| SBS1159+545 | 0.11 | 0.01 | 0.10 | 0.01 | 0.12 | 0.03 | 0.59 | 0.07 | $12^{d}$ |  |
| SBS1211+540 | 0.16 | 0.02 | 0.08 | 0.01 | $\leq 0.12$ | - | 0.30 | 0.05 | $15^{a}$ |  |

Table 6.9. IRAC flux densities for the DGS galaxies. When an upper limit is given, it is the $3 \sigma$ upper limit computed in Section 6.2.1 Photometry.

| Source | $\mathrm{F}_{3.6}$ [mJy] | $\sigma_{3.6}[\mathrm{mJy}]$ | $\mathrm{F}_{4.5}$ [mJy] | $\sigma_{4.5}[\mathrm{mJy}]$ | $\mathrm{F}_{5.8}$ [mJy] | $\sigma_{5.8}[\mathrm{mJy}]$ | $\mathrm{F}_{8.0}$ [mJy] | $\sigma_{8.0}[\mathrm{mJy}]$ | Radius ( ${ }^{\prime \prime}$ ) | Aper corr ? |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| SBS1249+493 | 0.08 | 0.02 | 0.07 | 0.02 | $\leq 0.20$ | - | 0.26 | 0.04 | $5^{\text {b }}$ |  |
| SBS1415+437 | - | - | 0.80 | 0.11 | - | - | 0.50 | 0.08 | $34^{a}$ | yes |
| SBS1533+574 | 2.09 | 0.34 | 1.52 | 0.24 | 2.26 | 0.33 | 3.32 | 0.48 | $30^{a}$ | yes |
| Tol0618-402 | 2.68 | 0.38 | 1.64 | 0.23 | 1.13 | 0.16 | 0.86 | 0.12 | $18^{a}$ | yes |
| Tol1214-277 | 0.08 | 0.01 | 0.09 | 0.01 | 0.15 | 0.03 | 0.25 | 0.03 | $8^{\text {b }}$ |  |
| UGC4483 | 1.69 | 0.25 | 1.02 | 0.15 | $\leq 0.33 / \leq 0.85^{g}$ | - | 0.90 | 0.15 | $43 \times 24(32)^{c}$ | yes |
| UGCA20 | $\leq 0.85$ | - | $\leq 0.60$ | - | $\leq 1.07$ | - | $\leq 0.46$ | - | $20^{a}$ |  |
| UM133 | $\leq 0.62$ | - | $\leq 0.46$ | - | $\leq 0.86$ | - | $\leq 0.54$ | - | $8^{a}$ |  |
| UM311 | - | - | 33.4 | 4.7 | - | - | 114 | 16 | $115^{e}$ | yes |
| UM448 | 17.7 | 2.5 | 13.2 | 1.9 | 34.1 | 4.8 | 83.0 | 11.7 | $53 \times 31(41)^{c}$ | yes |
| UM461 | 0.83 | 0.12 | 0.61 | 0.09 | 0.81 | 0.14 | 1.58 | 0.29 | $17^{a}$ | yes |
| VIIZw403 | 2.86 | 0.41 | 2.23 | 0.32 | 1.69 | 0.25 | 2.56 | 0.37 | $40^{a}$ | yes |

[^32]Table 6.10. WISE flux densities for the DGS galaxies.

| Source | $\mathrm{F}_{3.4}$ [mJy] | $\sigma_{3.4}[\mathrm{mJy}]$ | $\mathrm{F}_{4.6}$ [mJy] | $\sigma_{4.6}$ [mJy] | $\mathrm{F}_{12}$ [mJy] | $\sigma_{12}$ [mJy] | $\mathrm{F}_{22.0}$ [mJy] | $\sigma_{22.0}[\mathrm{mJy}$ ] | Extended source? |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Haro11 | 17.3 | 0.1 | 33.2 | 0.2 | 322 | 1 | 1957 | 12 | yes |
| Haro2 | 16.0 | 0.1 | 11.8 | 0.1 | 138 | 1 | 743 | 5 | yes |
| Haro3 | 12.5 | 0.1 | 9.8 | 0.1 | 130 | 1 | 739 | 5 | yes |
| He2-10 | 66.7 | 0.4 | 53.0 | 0.3 | 775 | 4 | 4696 | 29 | yes |
| HS0017+1055 | 0.148 | 0.001 | 0.291 | 0.002 | 3.54 | 0.02 | 12.1 | 0.1 |  |
| HS0052+2536 | 0.534 | 0.003 | 0.397 | 0.002 | 3.28 | 0.02 | 19.6 | 0.1 |  |
| HS0822+3542 | 0.106 | 0.001 | 0.0823 | 0.0004 | $\leq 0.47$ | - | 3.85 | 0.02 |  |
| HS1222+3741 | 0.125 | 0.001 | 0.122 | 0.001 | 1.08 | 0.01 | 5.94 | 0.04 |  |
| HS1236+3937 ${ }^{\text {a }}$ | - | - | - | - | - | - | - | - | - |
| HS1304+3529 | 0.221 | 0.001 | 0.221 | 0.001 | 1.43 | 0.01 | 10.5 | 0.1 |  |
| HS1319+3224 | 0.0392 | 0.0002 | 0.0801 | 0.0004 | 0.504 | 0.002 | 1.93 | 0.01 |  |
| HS1330+3651 | 0.274 | 0.002 | 0.212 | 0.001 | 1.09 | 0.01 | 7.64 | 0.05 |  |
| HS1442+4250 | 0.128 | 0.001 | 0.090 | 0.001 | $\leq 0.15$ | - | $\leq 1.05$ | - |  |
| HS2352+2733 | 0.0664 | 0.0004 | 0.0406 | 0.0002 | $\leq 0.42$ | - | $\leq 2.98$ | - |  |
| IZw18 | 0.212 | 0.001 | 0.207 | 0.001 | 0.773 | 0.004 | 5.44 | 0.03 |  |
| IC10 | 1204 | 7 | 769 | 4 | 2563 | 12 | 9088 | 93 | yes |
| IIZw40 | 6.29 | 0.03 | 11.5 | 0.1 | 289 | 1 | 1456 | 9 | yes |
| Mrk1089 | 5.74 | 0.03 | 5.18 | 0.03 | 69.0 | 0.3 | 412 | 3 | yes |
| Mrk1450 | 0.92 | 0.01 | 0.712 | 0.004 | 7.03 | 0.03 | 49.6 | 0.3 |  |
| Mrk153 | 1.64 | 0.01 | 1.03 | 0.01 | 3.92 | 0.02 | 25.4 | 0.2 | only at $3.4 \mu \mathrm{~m}$ |
| Mrk209 | 0.93 | 0.01 | 0.588 | 0.003 | 5.95 | 0.03 | 48.1 | 0.3 | only at $3.4 \mu \mathrm{~m}$ |
| Mrk930 | 1.63 | 0.01 | 1.41 | 0.01 | 15.3 | 0.1 | 114 | 1 | yes |
| NGC1140 | 15.0 | 0.1 | 10.2 | 0.1 | 57.3 | 0.3 | 310 | 2 | yes |
| NGC1569 ${ }^{\text {b }}$ | - | - | - | - | - | - | - | - | - |
| NGC1705 | 9.89 | 0.05 | 6.42 | 0.04 | 8.16 | 0.04 | 30.7 | 0.2 | yes |
| NGC2366 ${ }^{\text {c }}$ | - | - | - | - | - | - | - | - | - |
| NGC4214 | 10.9 | 0.1 | 8.94 | 0.05 | $86.7^{e}$ | 0.4 | $689{ }^{\text {e }}$ | 4 | yes |
| NGC4449 | 389 | 2 | 249 | 1 | 905 | 4 | 255 | 16 | yes |
| NGC4861 ${ }^{\text {b }}$ | - | - | - | - | - | - | - | - | - |
| NGC5253 | 62.0 | 0.3 | 168 | 1 | 1391 | 6 | 6562 | 58 | yes |
| NGC625 ${ }^{\text {c }}$ | 7.24 | 0.04 | 5.01 | 0.03 | 35.1 | 0.2 | 380 | 2 | yes |
| NGC6822 ${ }^{\text {b }}$ | - | - | - | - | - | - | - | - | - |
| Pox186 | 0.106 | 0.001 | 0.087 | 0.001 | 2.18 | 0.01 | 13.3 | 0.1 |  |
| SBS0335-052 | 0.460 | 0.003 | 1.67 | 0.01 | 24.4 | 0.1 | 74.1 | 0.5 |  |
| SBS1159+545 | 0.0652 | 0.0004 | 0.095 | 0.001 | 1.00 | 0.01 | 7.18 | 0.04 |  |
| SBS1211+540 | 0.094 | 0.001 | 0.0642 | 0.0003 | 0.225 | 0.001 | 1.86 | 0.01 |  |

Table 6.10. WISE flux densities for the DGS galaxies.

| Source | $\mathrm{F}_{3.4}$ [mJy] | $\sigma_{3.4}$ [mJy] | $\mathrm{F}_{4.6}$ [mJy] | $\sigma_{4.6}[\mathrm{mJy}]$ | $\mathrm{F}_{12}$ [mJy] | $\sigma_{12}$ [mJy] | $\mathrm{F}_{22.0}$ [mJy] | $\sigma_{22.0}$ [mJy] | Extended source? |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| SBS1249+493 | 0.0507 | 0.0003 | 0.0684 | 0.0004 | 0.726 | 0.003 | 3.40 | 0.02 |  |
| SBS1415+437 | 0.283 | 0.002 | 0.286 | 0.002 | 1.75 | 0.01 | 17.9 | 0.1 |  |
| SBS1533+574 | 0.96 | 0.01 | 0.779 | 0.004 | 8.04 | 0.04 | 49.7 | 0.3 | only at $3.4 \mu \mathrm{~m}$ |
| Tol0618-402 | 1.70 | 0.01 | 0.92 | 0.01 | 0.281 | 0.001 | $\leq 1.00$ | - |  |
| Tol1214-277 | 0.0513 | 0.0003 | 0.0786 | 0.0004 | 1.24 | 0.01 | 6.89 | 0.04 |  |
| UGC4483 ${ }^{\text {c }}$ | - | - | - | - | - | - | - | - | - |
| UGCA20 | 0.0667 | 0.0004 | 0.0346 | 0.0002 | $\leq 0.19$ | - | $\leq 1.30$ | - |  |
| UM133 | 0.209 | 0.001 | 0.171 | 0.001 | 0.695 | 0.003 | 5.75 | 0.04 |  |
| UM311 ${ }^{\text {d }}$ | - | - | - | - | - | - | - | - | - |
| UM448 | 11.8 | 0.1 | 9.74 | 0.05 | 93.5 | 0.4 | 543 | 3 | yes |
| UM461 | 0.346 | 0.002 | 0.372 | 0.002 | 5.51 | 0.03 | 29.2 | 0.2 |  |
| VIIZw403 ${ }^{\text {c }}$ | - | - | - | - | - | - | - | - | - |

[^33]Table 6.11. IRAS flux densities for the DGS galaxies.

| Source | $\mathrm{F}_{12}$ [Jy] | $\sigma_{F_{12}}$ [Jy] | $\mathrm{F}_{25}$ [Jy] | $\sigma_{F_{25}}[\mathrm{Jy}]$ | $\mathrm{F}_{60}[\mathrm{Jy}]$ | $\sigma_{F_{60}}[\mathrm{Jy}]$ | $\mathrm{F}_{100}$ [Jy] | $\sigma_{F_{100}}[\mathrm{Jy}]$ | Ref |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Haro11 | 0.42 | 0.05 | 2.49 | 0.03 | 6.48 | 0.39 | 5.01 | 0.35 | 1 |
| Haro2 | 0.21 | 0.03 | 0.95 | 0.01 | 4.68 | 0.28 | 5.32 | 0.32 | 1 |
| Haro3 | 0.21 | 0.03 | 0.94 | 0.01 | 4.95 | 0.40 | 6.75 | 0.41 | 1 |
| He2-10 | 1.18 | 0.13 | 6.78 | 0.72 | 23.4 | 2.9 | 26.3 | 3.9 | 2 |
| HS0017+1055 | - | - | - | - | - | - | - | - | - |
| HS0052+2536 | $0.11{ }^{a}$ | 0.02 | $0.14{ }^{a}$ | 0.01 | $0.25{ }^{\text {a }}$ | 0.01 | $0.71{ }^{\text {a }}$ | 0.07 | 1 |
| HS0822+3542 | - | - | - | - | - | - | - | - | - |
| HS1222+3741 | - | - | - | - | - | - | - | - | - |
| HS1236+3937 | - | - | - | - | - | - | - | - | - |
| HS1304+3529 | - | - | - | - | - | - | - | - | - |
| HS1319+3224 | - | - | - | - | - | - | - | - | - |
| HS1330+3651 | - | - | - | - | - | - | - | - | - |
| HS1442+4250 | - | - | - | - | - | - | - | - | - |
| HS $2352+2733$ | - | - | - | - | - | - | - | - | - |
| IZw18 | - | - | - | - | - | - | - | - | - |
| IC10 | - | - | - | - | - | - | - | - | - |
| IIZw40 | 0.41 | 0.04 | 1.88 | 0.02 | 6.02 | 0.42 | $\leq 19.7$ | - | 1 |
| Mrk1089 | $0.25{ }^{\text {a }}$ | 0.01 | 0.70 | 0.01 | 4.06 | 0.20 | 5.64 | 0.56 | 3 |
| Mrk1450 | $\leq 0.06$ | - | $\leq 0.10$ | - | 0.28 | 0.04 | $\leq 0.58$ | - | 1 |
| Mrk153 | $\leq 0.08$ | - | $\leq 0.09$ | - | 0.28 | 0.04 | $\leq 0.48$ | - | 1 |
| Mrk209 | - | - | - | - | - | - | - | - | - |
| Mrk930 | $\leq 0.08$ | - | 0.23 | 0.01 | 1.25 | 0.09 | $\leq 2.15$ | - | 1 |
| NGC1140 | 0.10 | 0.03 | 0.49 | 0.06 | 3.34 | 0.31 | 4.92 | 0.64 | 2 |
| NGC1569 | 0.87 | 0.07 | 7.73 | 0.62 | 44.0 | 3.5 | 47.1 | 5.9 | 2 |
| NGC1705 | $\leq 0.05$ | - | $\leq 0.11$ | - | $\leq 0.87$ | - | $\leq 1.82$ | - | 1 |
| NGC2366 | $\leq 0.12$ | - | 0.70 | 0.07 | $3.51{ }^{\text {a }}$ | 0.18 | $4.67{ }^{\text {a }}$ | 0.28 | 1 |
| NGC4214 | 0.65 | 0.07 | 2.58 | 0.19 | 17.9 | 1.2 | 29.0 | 3.2 | 2 |
| NGC4449 | $2.14{ }^{\text {a }}$ | 0.23 | 5.15 | 0.65 | 36.6 | 4.0 | 73.0 | 11.3 | 2 |
| NGC4861 | $0.25{ }^{\text {a }}$ | 0.03 | 0.41 | 0.09 | 1.82 | 0.38 | 2.39 | 0.55 | 2 |
| NGC5253 | $2.81{ }^{\text {a }}$ | 0.20 | 12.3 | 1.0 | 29.0 | 2.3 | 29.1 | 3.6 | 2 |
| NGC625 | 0.20 | 0.03 | 1.30 | 0.03 | 5.73 | 0.04 | 8.63 | 0.13 | 4 |
| NGC6822 | $0.25{ }^{\text {a }}$ | 0.04 | 2.46 | 0.37 | 47.6 | 7.1 | 95.4 | 14.3 | 5 |
| Pox186 | - | - | - | - | - | - | - | - | - |
| SBS0335-052 | - | - | - | - | - | - | - | - | - |
| SBS1159+545 | - | - | - | - | - | - | - | - | - |
| SBS1211+540 | - | - | - | - | - | - | - | - | - |
| SBS1249+493 | - | - | - | - | - | - | - | - | - |
| SBS1415+437 | - | - | - | - | - | - | - | - | - |
| SBS1533+574 | $\leq 0.06$ | - | $\leq 0.07$ | - | $0.26^{a}$ | 0.03 | $0.41{ }^{a}$ | 0.11 | 1 |
| Tol0618-402 | - | - | - | - | - | - | - | - | - |
| Tol1214-277 | - | - | - | - | - | - | - | - | - |
| UGC4483 | $\leq 0.09$ | - | $\leq 0.09$ | - | $\leq 0.12$ | - | $\leq 0.57$ | - | 2 |
| UGCA20 |  | - | - | - | - | - | - | - | - |
| UM133 | - | - | - | - | - | - | - | - | - |
| UM311 | - | - | - | - | - | - | - | - | - |
| UM448 | $0.15{ }^{\text {a }}$ | 0.04 | 0.78 | 0.11 | 4.01 | 0.40 | 4.30 | 0.61 | 2 |
| UM461 | $\leq 0.09$ | - | $\leq 0.09$ | - | $\leq 0.12^{a}$ | - | $\leq 0.57$ | - | 2 |
| VIIZw403 | $0.07^{a}$ | 0.01 | 0.06 | 0.01 | 0.39 | 0.04 | $0.90^{a}$ | 0.20 | 2 |

${ }^{a}$ : This flux density is not consistent with the rest of the IR photometry. Thus we do not consider it for the modelling.
References: (1) the NASA/IPAC ISA IRAS Faint Source Catalog (v2.0) available at: http://irsa.ipac.caltech.edu/cgi-bin/Gator/nph-dd?catalog=irasfsc ; (2) Engelbracht et al. (2008) ; (3) the NASA/IPAC ISA IRAS Point Source Catalog (v2.1) available at: http://irsa.ipac.caltech.edu/cgi-bin/Gator/nph-dd?catalog=iraspsc ; (4) Sanders et al. (2003) ; (5) Rice et al. (1988)

Table 6.12. Submm flux densities for the DGS sample.

| Name | $\mathrm{F}_{450} \mu \mathrm{~m}[\mathrm{mJy}]$ | Ref | $\mathrm{F}_{850} \mu \mathrm{~m}$ [mJy] | Ref | $\mathrm{F}_{870} \mu \mathrm{~m}$ [mJy] | Ref | $\mathrm{F}_{1200} \mu \mathrm{~m}$ [mJy] | Ref |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Haro11 |  | - | - | - | $44.0 \pm 6.0$ | 1 | - | - |
| - Radio | - | - | - | - | $12.2 \pm 8.2$ | 2,3 | - | - |
| - $\mathrm{CO}(3-2)$ | - | - | - | - | $0.3 \pm 0.1$ | $4^{a}$ | - | - |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | - | - |
| $\mathrm{F}_{\text {dust }}$ | - | - | - | - | $31.5 \pm 10.2$ | - | - | - |
| Haro2 | - | - | $40.0 \pm 10.0$ | 5 | - | - | - | - |
| - Radio | - | - | $9.8 \pm 1.2$ | 5 | - | - | - | - |
| - $\mathrm{CO}(3-2)$ | - | - | $3.2 \pm 0.7$ | $6^{a}$ | - | - | - | - |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | - | - |
| $\mathrm{F}_{\text {dust }}$ | - | - | $27.0 \pm 10.1$ | - | - | - | - | - |
| He2-10 | $342 \pm 65$ | 7 | $130 \pm 12$ | 7 | - | - | $60 \pm 14$ | 7 |
| - Radio | - | - | $20.0 \pm 6.7$ | 8 | - | - | $20.7 \pm 7.0$ | 8 |
| - $\mathrm{CO}(3-2)$ | - | - | $3.0 \pm 0.0$ | $7,9^{a}$ | - | - | - | - |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | $5.0 \pm 0.0$ | $7,10^{\text {b }}$ |
| $\mathrm{F}_{\text {dust }}$ | $342 \pm 65$ | - | $107 \pm 14$ | - | - | - | $34.3 \pm 15.6$ | - |
| IIZw40 | $248 \pm 81$ | 7 | $98 \pm 14$ | 7 | - | - | $43.0 \pm 13.0$ | 7 |
| - Radio | - | - | $13.5 \pm 4.5$ | 3,11 | - | - | $14.0 \pm 4.7$ | 3,11 |
| - $\mathrm{CO}(3-2)$ | - | - | $\leq 3.4$ | $7,9^{a}$ | - | - | - |  |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | $0.1 \pm 0.0$ | $7,12^{\text {b }}$ |
| $\mathrm{F}_{\text {dust }}$ | $248 \pm 81$ | - | $81.1 \pm 14.7$ | - | - | - | $28.9 \pm 13.8$ | - |
| Mrk1089 | - | - | - | - | $74.8 \pm 10.1$ | 1 | - | - |
| - Radio | - | - | - | - | $4.6 \pm 0.0$ | 11 | - | - |
| - $\mathrm{CO}(3-2)$ | - | - | - | - | $0.6 \pm 0.2$ | $4^{a}$ | - | - |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | - | - |
| $\mathrm{F}_{\text {dust }}$ | - | - | - | - | $69.5 \pm 10.1$ | - | - | - |
| Mrk930 | - | - | - | - | $41.0 \pm 5.0$ | 13 | - | - |
| - Radio | - | - | - | - | $7.0 \pm 0.5$ | 3 | - | - |
| - $\mathrm{CO}(3-2)$ | - | - | - | - | $\leq 0.03$ | $4^{\text {b }}$ | - | - |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | - | - |
| $\mathrm{F}_{\text {dust }}$ | - | - | - | - | $33.9 \pm 5.0$ | - | - | - |
| NGC1140 | $272 \pm 55$ | 7 | $69.0 \pm 28.0$ | 7 | - | - | $\leq 48.0$ | 7 |
| - Radio | - | - | $7.8 \pm 0.0$ | 14 | - | - | $8.1 \pm 0.0$ | 14 |
| - $\mathrm{CO}(3-2)$ | - | - | $\leq 0.01$ | $15^{b}$ | - | - | - | - |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | $\leq 0.02$ | $15^{b}$ |
| $\mathrm{F}_{\text {dust }}$ | $272 \pm 55$ | - | $61.2 \pm 28.0$ | - | - | - | $\leq 39.9$ | - |
| NGC1569 | $1280 \pm 450$ | 16 | $345 \pm 40$ | 16 | - | - | $250 \pm 99$ | 17 |
| - Radio ${ }^{\text {e }}$ | - | - | $49.0 \pm 0.0$ | 16 | - | - | $55.5 \pm 0.0$ | 16 |
| - $\mathrm{CO}(3-2)$ | - | - | $1.7 \pm 0.2$ | $9^{a}$ | - | - | - |  |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | $0.11 \pm 0.01$ | $18^{b}$ |
| $\mathrm{F}_{\text {dust }}$ | $1280 \pm 450$ | - | $294 \pm 40$ | - | - | - | $194 \pm 99$ | - |
| NGC1705 | - | - | - | - | $127 \pm 17$ | 1 | - | - |
| - Radio | - | - | - | - | $7.0 \pm 0.5$ | 19 | - | - |
| - $\mathrm{CO}(3-2)$ | - | - | - | - | $1.0 \pm 0.0$ | - ${ }^{d}$ | - | - |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | - | - |
| $\mathrm{F}_{\text {dust }}$ | - | - | - | - | $119 \pm 17$ | - | - | - |
| NGC4214 | - | - | - | - | - | - | $350 \pm 106$ | 20 |
| - Radio | - | - | - | - | - | - | $23.6 \pm 6.7$ | 21 |
| - $\mathrm{CO}(3-2)$ | - | - | - | - | - | - | - | - |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | $0.12 \pm 0.02$ | $22^{\text {c }}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | - | - | - | - | $326 \pm 106$ | - |
| NGC5253 | - | - | - | - | $494 \pm 59$ | 13 | - | - |
| - Radio | - | - | - | - | $50.4 \pm 2.0$ | 3 | - | - |
| - $\mathrm{CO}(3-2)$ | - | - | - | - | $\leq 0.4$ | $23^{a}$ | - | - |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | - | - |
| $\mathrm{F}_{\text {dust }}$ | - | - | - | - | $443 \pm 59$ | - | - | - |
| UM448 | - | - | - | - | $76.0 \pm 9.0$ | 13 | - | - |
| - Radio | - | - | - | - | $18.8 \pm 0.8$ | 3 | - | - |
| - $\mathrm{CO}(3-2)$ | - | - | - | - | $1.9 \pm 0.6$ | $12^{\text {c }}$ | - | - |
| - $\mathrm{CO}(2-1)$ | - | - | - | - | - | - | - | - |
| $\mathrm{F}_{\text {dust }}$ | - | - | - | - | $55.4 \pm 9.1$ | - | - | - |

${ }^{a}$ : From CO(3-2).
${ }^{b}$ : From $\mathrm{CO}(2-1)$.
${ }^{c}$ : From CO(1-0).
${ }^{d}$ : Mean line contamination.
${ }^{e}$ : For this galaxy, Galliano et al. (2003) used a spectral index for the radio contamination $\alpha_{\text {radio }}=-0.36$ determined by Israel \& de Bruyn (1988) from the full radio spectrum of NGC1569.
Note : When the error is 0.0 , it means that no error was given on the original radio or CO measurement.
References : (1) Galametz et al. (2009) ; (2) Mauch et al. (2008) ; (3) Condon et al. (1998) ; (4) Cormier et al. (201456 (5) Dale et al. (2007) ; (6) Wilson et al. (2012); (7) Galliano et al. (2005); (8) Kobulnicky \& Johnson (1999); (9) Meier et al. (2001); (10) Baas et al. (1994); (11) Sramek \& Weedman (1986); (12) Sage et al. (1992); (13) this work; (14) Klein et al. (1983); (15) Hunter \& Sage (1993); (16) Galliano et al. (2003); (17) Lisenfeld et al. (2002); (18) Greve et al. (1996); (19) Meurer et al. (1998); (20) Hermelo et al. (2013); (21) Kepley et al. (2011); (22) Walter et al. (2001); (23) Hirashita (2013)

Table 6.13. Submm flux densities for the KINGFISH sample.

| Name | $\mathrm{F}_{450} \mu \mathrm{~m}[\mathrm{Jy}]$ | Ref | $\mathrm{F}_{850} \mu \mathrm{~m}$ [mJy] | Ref |
| :---: | :---: | :---: | :---: | :---: |
| NGC0337 | - | - | $350 \pm 50$ | 1 |
| - Radio | - | - | $63.7 \pm 6.4$ | 1 |
| - CO(3-2) | - | - | $\leq 6.4$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $280 \pm 50$ | - |
| NGC1097 | - | - | $1440 \pm 780$ | 1 |
| - Radio | - | - | $240 \pm 24$ | 1 |
| - CO(3-2) | - | - | $108 \pm 1$ | $3^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $1092 \pm 780$ | - |
| NGC1482 | - | - | $330 \pm 50$ | 1 |
| - Radio | - | - | $138 \pm 14$ | 1 |
| - CO(3-2) | - | - | $2.8 \pm 0.3$ | $4^{b}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $189 \pm 52$ | - |
| NGC2798 | - | - | $190 \pm 30$ | 1 |
| - Radio | - | - | $48.1 \pm 5.2$ | 1 |
| - $\mathrm{CO}(3-2)$ | - | - | $20.2 \pm 2.2$ | $5^{b}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $122 \pm 31$ | - |
| NGC2976 | - | - | $610 \pm 240$ | 1 |
| - Radio | - | - | $29.5 \pm 2.9$ | 1 |
| - $\mathrm{CO}(3-2)$ | - | - | $23.4 \pm 3.7$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $557 \pm 240$ | - |
| NGC3190 | - | - | $190 \pm 40$ | 1 |
| - Radio | - | - | $24.9 \pm 2.9$ | 1 |
| - $\mathrm{CO}(3-2)$ | - | - | $\leq 3.6$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $162 \pm 40$ | - |
| NGC3521 | - | - | $2110 \pm 820$ | 1 |
| - Radio | - | - | $207 \pm 21$ | 1 |
| - $\mathrm{CO}(3-2)$ | - | - | $174 \pm 12$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $1729 \pm 820$ | - |
| NGC3627 | - | - | $1860 \pm 700$ | 1 |
| - Radio | - | - | $265 \pm 27$ | 1 |
| - $\mathrm{CO}(3-2)$ | - | - | $303 \pm 17$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $1292 \pm 701$ | - |
| NGC4254 | - | - | $1010 \pm 540$ | 1 |
| - Radio | - | - | $245 \pm 24$ | 1 |
| - $\mathrm{CO}(3-2)$ | - | - | $161 \pm 10$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $604 \pm 541$ | - |
| NGC4321 | - | - | $880 \pm 490$ | 1 |
| - Radio | - | - | $197 \pm 20$ | 1 |
| - $\mathrm{CO}(3-2)$ | - | - | $119 \pm 12$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $564 \pm 491$ | - |
| NGC4536 | - | - | $420 \pm 110$ | 1 |
| - Radio | - | - | $112 \pm 11$ | 1 |
| - CO (3-2) | - | - | $20.7 \pm 2.3$ | $6^{c}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $287 \pm 111$ | - |
| NGC4569 | - | - | $470 \pm 80$ | 1 |
| - Radio | - | - | $48.1 \pm 5.2$ | 1 |
| - CO(3-2) | - | - | $43.8 \pm 3.7$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $378 \pm 80$ | - |
| NGC4579 | - | - | $440 \pm 70$ | 1 |
| - Radio | - | - | $56.8 \pm 5.8$ | 1 |
| - CO(3-2) | - | - | $17.2 \pm 3.0$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $366 \pm 70$ | - |
| NGC4594 | - | - | $370 \pm 110$ | 1 |
| - Radio | - | - | $79.4 \pm 8.1$ | 1 |
| - CO(3-2) | - | - | $\leq 4.4$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $286 \pm 110$ | - |
| NGC4631 | $30.7 \pm 10.0$ | 1 | $5730 \pm 1210$ | 1 |
| - Radio | - | - | $695 \pm 70$ | 1 |
| - $\mathrm{CO}(3-2)$ | - | - | $150 \pm 7$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | $30.7 \pm 10.0$ | - | $4885 \pm 1212$ | - |
| NGC4736 | - | 1 | $1540 \pm 660$ | 1 |
| - Radio | - | - | $157 \pm 16$ | 1 |
| - CO(3-2) | - | - | $117 \pm 9$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $1266 \pm 660$ | - |
| NGC4826 | - | 1 | $1230 \pm 310$ | 1 |
| - Radio | - | - | $58.5 \pm 5.8$ | 1 |
| - CO(3-2) | - | - | $218 \pm 11$ | $2^{a}$ |
| $\mathrm{F}_{\text {dust }}$ | - | - | $954 \pm 310$ | - |

Table 6.13. Submm flux densities for the KINGFISH sample.

| Name | $\mathrm{F}_{450} \mu \mathrm{~m}[\mathrm{Jy}]$ | Ref | $\mathrm{F}_{850} \mu \mathrm{~m}[\mathrm{mJy}]$ | Ref |
| :--- | :---: | :---: | :---: | :---: |
| NGC5713 | - | - | $570 \pm 120$ | 1 |
| - Radio | - | - | $92.7 \pm 9.3$ | 1 |
| - CO(3-2) | - | - | $6.4 \pm 0.7$ | $4^{b}$ |
| $\mathrm{~F}_{\text {dust }}$ | - | - | $471 \pm 120$ | - |
| NGC5866 | $0.8 \pm 0.2$ | 1 | $140 \pm 20$ | 1 |
| - Radio | - | - | $13.3 \pm 1.7$ | 1 |
| - CO(3-2) | - | - | $14.1 \pm 1.7$ | $7^{c}$ |
| $\mathrm{~F}_{\text {dust }}$ | $0.8 \pm 0.2$ | - | $113 \pm 20$ | - |
| NGC6946 | $18.5 \pm 4.6$ | 1 | $2980 \pm 450$ | 1 |
| - Radio | - | - | $808 \pm 81$ | 1 |
| - CO(3-2) | - | - | $553 \pm 61$ | $5^{b}$ |
| $\mathrm{~F}_{\text {dust }}$ | $18.5 \pm 4.6$ | - | $1619 \pm 461$ | - |
| NGC7331 | $20.6 \pm 8.1$ | 1 | $2110 \pm 380$ | 1 |
| - Radio | - | - | $216 \pm 21$ | 1 |
| - CO(3-2) $^{\mathrm{F}_{\text {dust }}}$ | - | - | $114 \pm 13$ | $5^{b}$ |

${ }^{a}$ : From $\mathrm{CO}(3-2)$.
${ }^{b}$ : From $\mathrm{CO}(2-1)$.
${ }^{c}$ : From CO(1-0).
References : (1) Dale et al. (2007) ; (2) Wilson et al. (2012) ; (3) Petitpas \& Wilson (2003) ; (4) Albrecht et al. (2007) ; (5) Leroy et al. (2009) ; (6) Schruba et al. (2012); (7) Welch \& Sage (2003)

## Chapter 7

## Modelling the dust emission of galaxies

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In this Chapter, we present the semi-empirical SED model that will enable us to derive more accurate estimates of the dust properties (especially the dust masses) compared to those determined using modified blackbodies in Chapter 5 (Section 7.1). In Section 7.2, we then describe how this model is applied it to the DGS and KINGFISH galaxies.

### 7.1 Description of the model

### 7.1.1 Philosophy of the model

As we already understood from Chapter 1, modelling the dust emission of a galaxy is a very complex process. Several regions of the galaxy such as star forming regions, molecular clouds, diffuse ISM, etc. contribute differently to the dust SED at each wavelength, and all of these contributions can be captured in the telescope beam. The illumination conditions and the dust composition can vary between the different regions or within the region itself. However, due to lack of observational constraints and spatial resolution we can not access this level of detail in galaxies at the longest wavelengths, except perhaps for the closest galaxies (i.e., LMC/SMC: D~50 kpc, or M31: D~800 $\mathrm{kpc})$. We thus model here the global emission of our galaxies.

In a given region within a galaxy the dust physical conditions are directly linked to the illumination conditions. For example, dust within molecular clouds, where little radiation can penetrate, is colder than dust in the diffuse ISM. In the model, we suppose that each region of the galaxy is subject to uniform illumination conditions and that the dust composition and properties are the same throughout the galaxy. We account for the different regions in a galaxy by combining different regions with different interstellar radiation field intensities. The model has thus two steps:

1. First, we model the dust SED of a mass-element of the ISM, characterised by a uniform intensity of the radiation field, and controlled by the microscopic grain properties.
2. Then, we combine several mass-elements to account for the wide variety of illumination conditions within the galaxy.

This model is a simplified version of the model from Galliano et al. (2008), and is presented in Galliano et al. (2011). It has been used to model galaxies before by Galametz et al. (2009, 2010, 2011); Hony et al. (2010); Meixner et al. (2010); O’Halloran et al. (2010).

### 7.1.2 Grain properties

To build the dust model, we require a set of dust components. Each dust specie is characterised by a chemical composition and a size distribution function. The chemical composition of the grains then determines the set of optical properties that is used. We adopt a standard dust composition from Zubko et al. (2004): the Galactic grain composition made of 3 dust species: PAHs, graphite and silicate grains (i.e., the BARE-GR-S model from Zubko et al. 2004). To determine the contribution to the total emission of each dust specie, one needs to know the size distribution function and the optical properties of each dust component.

## Size distribution

The size distribution of each dust component that we use in our modelling has been determined by Zubko et al. (2004) and was constrained by fitting simultaneously the UV extinction curve, the IR emission and the elemental depletions of the diffuse Galactic ISM. We outline here the method followed by Zubko et al. (2004).

Let $F_{i}{ }^{1}$ be the size distribution of the $\mathrm{i}^{\text {th }}$ dust specie with $\mathrm{i}=\{1,2,3\}=\{\mathrm{PAH}$, graphite, silicate $\}$ and $F_{i}(a) d a$ the number of grains per H atom in the radius range $[\mathrm{a}, \mathrm{a}+\mathrm{da}]$. Let $k$ be the number of chemical elements constituting the dust. Observations of the diffuse Galactic ISM give the following set of constraints, used by Zubko et al. (2004):

1. Constraints from the abundance of elements locked up into the solid phase of the ISM:

$$
\begin{equation*}
\frac{N_{j}}{N_{H}}=\sum_{i=1}^{3} \int \frac{4}{3} \pi a^{3} \rho_{i} \frac{\alpha_{i j}}{m_{j}} F_{i}(a) \mathrm{d} a \tag{7.1}
\end{equation*}
$$

where $N_{j}$ is the column density of the $\mathrm{j}^{\text {th }}$ chemical element constituting the dust specie ( $\mathrm{j}=$ $1 \ldots \mathrm{k}$ ), locked up in the dust; $N_{H}$ is the line-of-sight hydrogen column density; $\rho_{i}$ is the mass density of the $\mathrm{i}^{\text {th }}$ dust component; $m_{j}$ is the atomic mass of the $\mathrm{j}^{\text {th }}$ chemical element; and $\alpha_{i j}$ is the mass fraction of the $\mathrm{j}^{\text {th }}$ element in the $\mathrm{i}^{\text {th }}$ dust specie.
Input from the observations of the diffuse Galactic ISM: The relative abundance of an element locked up in dust, $N_{j} / N_{H}$, is given by the elemental depletion in the gas phase: the difference between the total ISM abundance of the elements and its observed gas phase abundance. The first is determined via the combination of stellar surface abundances (solar, B, F and G types) from Snow \& Witt (1996); Holweger (2001); Sofia \& Meyer (2001). The observed gas phase abundances are from Cardelli et al. (1996); Dwek et al. (1997); Meyer et al. (1997, 1998). The densities, $\rho_{i}$, are given in Table 7.1, and the mass fractions, $\alpha_{i j}$, can be calculated from the atomic and molecular masses of the dust constituents.
2. Constraints from the UV-visible interstellar extinction curve:

$$
\begin{equation*}
\frac{\tau_{H}(\lambda)}{N_{H}}=\sum_{i=1}^{3} \int \pi a^{2} Q_{e x t}^{i}(\lambda, a) F_{i}(a) \mathrm{d} a \tag{7.2}
\end{equation*}
$$

where $\tau_{H}(\lambda)$ is the extinction per H column density; and $Q_{\text {ext }}(a, \lambda)$ is the extinction efficiency of the $\mathrm{i}^{\text {th }}$ dust component at radius $a$. Eq. 7.2 is similar to Eq. 1.35 from Section 1.2.3.
Input from the observations of the diffuse Galactic ISM: To characterise the extinction from the diffuse ISM, Zubko et al. (2004) use the mean extinction curve for $R_{V}=3.1$ of Fitzpatrick (1999). The extinction cross-section is calculated using Mie theory for spherical dust particles (Bohren \& Huffman 1983) and the dielectric function for the dust component. The references for the dielectric functions are listed in Table 7.1.
3. Constraints from the IR emission spectrum:

$$
\begin{equation*}
\frac{I_{\lambda}(\lambda)}{N_{H}}=\sum_{i=1}^{3} \int \pi a^{2} Q_{a b s}^{i}(\lambda, a) S_{\lambda}^{i}(\lambda, a) F_{i}(a) \mathrm{d} a \tag{7.3}
\end{equation*}
$$

where $I_{\lambda}(\lambda)$ is the specific intensity per unit solid angle; $Q_{a b s}^{i}(a, \lambda)$ is the absorption efficiency of the $\mathrm{i}^{\text {th }}$ dust component at radius $a$; and $S_{\lambda}^{i}(\lambda, a)$ is the corresponding source function ${ }^{2}$.

[^34]For all dust particles, we have:

$$
\begin{equation*}
S_{\lambda}^{i}(\lambda, a)=\int_{T_{C M B}}^{T_{s u b}} B_{\lambda}(\lambda, T) P(a, T) \mathrm{d} T \tag{7.4}
\end{equation*}
$$

where $T$ is the dust grain temperature and can range between the CMB radiation temperature, $T_{C M B}$, and the grain sublimation temperature, $T_{\text {sub }}$; and $P(a, T)$ is the temperature probability distribution defined in Chapter 1, Section 1.2.4 (see Eqs. 1.33 and 1.34).
This is where we link the emission of the dust particles to the ISRF to which they are exposed: dust is heated by the ISRF and the heating depends on the grain size (see Section 1.2.3). Small grains undergo stochastic heating and present temperature and internal energy fluctuations depending on the flux of photons they receive. Big grains are in thermodynamical equilibrium with the ISRF and have an equilibrium temperature $T_{d}(a)$, i.e., $P(a, T)$ can be approximated by a delta function at $T_{d}(a)$, corresponding to the equilibrium temperature. In this case, $S_{\lambda}^{i}(\lambda, a) \simeq \mathrm{B}_{\lambda}\left(\lambda, T_{d}(a)\right)$.
As we saw in Chapter 1, $T_{d}(a)$ is controlled by the balance between what the dust particle absorbs and re-emits:

$$
\begin{equation*}
\int_{0}^{\infty} \pi a^{2} Q_{a b s}(\lambda, a) J_{\lambda}^{G a l}(\lambda) \mathrm{d} \lambda=\int_{0}^{\infty} \pi a^{2} Q_{a b s}(\lambda, a) B_{\lambda}\left(\lambda, T_{d}(a)\right) \mathrm{d} \lambda \tag{7.5}
\end{equation*}
$$

where $J_{\lambda}^{\text {Gal }}$ is the mean intensity of the ISRF in the solar neighbourhood.
Input from the observations of the diffuse Galactic ISM: IR data from the $\mathrm{COBE}^{3}$ all-sky survey in 8 DIRBE bands and in the FIRAS ${ }^{4}$ channels (Dwek et al. 1997; Arendt et al. 1998), is used to constrain the IR emission spectrum of the diffuse Galactic ISM, $I_{\lambda}(\lambda)$. The absorption cross-sections are computed the same way as the extinction cross-sections. The source functions, $S_{\lambda}^{i}(\lambda, a)$, are computed by Zubko et al. (2004) using the Thermal Discrete Approximation (TDA, as in Draine \& Li 2001) for small grains and is the Planck function at $T_{d}(a)$ for big grains. For the ISRF, Zubko et al. (2004) adopt the empirical representation for the solar neighbourhood from Mathis et al. (1983) with a mean intensity, $J_{\lambda}^{\text {Gal }}$ :

$$
\begin{equation*}
J_{\lambda}^{G a l}(\lambda)=J_{\lambda}^{U V}(\lambda)+\sum_{i=1}^{3} W_{i} B_{\lambda}\left(\lambda, T_{i}\right)+B_{\lambda}\left(\lambda, T_{C M B}\right) \tag{7.6}
\end{equation*}
$$

where $J_{\lambda}^{U V}$ is the UV component, the second term represents three effective blackbody sources with dilution factors $W_{\{1,2,3\}}=\left\{10^{-14}, 10^{-13}, 4 \times 10^{-13}\right\}$ and temperatures $T_{\{1,2,3\}}=\{7500$, $4000,3000\} \mathrm{K}$; and the third term represents the cosmic microwave background (CMB) radiation, $B_{\lambda}\left(\lambda, T_{C M B}\right)$ at the temperature $T_{C M B}=2.73 \mathrm{~K}$ (Mather et al. 1994).

The set of constraints used by Zubko et al. (2004) is presented in Fig. 7.1. Zubko et al. (2004) use the method of regularization to solve this set of constraints and obtain a size distribution for the chosen dust composition coherent with the observations.

For each dust specie, the size distribution, $F_{i}(a)$, is expressed as $F_{i}(a)=A \times G_{i}(a)$, where A is a normalisation coefficient in $\mathrm{H}^{-1}, a$, the grain radius, is expressed in $\mu \mathrm{m}$ and $G_{i}(a)$ is in units of $\mu \mathrm{m}^{-1}$ and follows:

[^35]\[

$$
\begin{equation*}
\int_{a_{i, \min }}^{a_{i, \max }} G_{i}(a) \mathrm{d} a=1 \tag{7.7}
\end{equation*}
$$

\]

where $a_{i, \text { min }}$ and $a_{i, \max }$ are the minimum and maximum grain size for the $\mathrm{i}^{\text {th }}$ dust specie (given in Table 7.1). An analytical function $\bar{G}$ is used to approximate the numerical values of $G(a)$ :

$$
\begin{equation*}
\log (\bar{G}(a))=c_{0}+b_{0} \log (a)-b_{1}\left|\log \left(\frac{a}{a_{1}}\right)\right|^{m_{1}}-b_{2}\left|\log \left(\frac{a}{a_{2}}\right)\right|^{m_{2}}-b_{3}\left|a-a_{3}\right|^{m_{3}}-b_{4}\left|a-a_{4}\right|^{m_{4}} \tag{7.8}
\end{equation*}
$$

The 14 parameters $\left\{c_{0}, b_{0}, b_{1}, b_{2}, b_{3}, b_{4}, a_{0}, a_{1}, a_{2}, a_{3}, a_{4}, m_{1}, m_{2}, m_{3}, m_{4}\right\}$ are given in Zubko et al. (2004).

The resulting size distribution is presented in Fig. 7.2.


Fig. 7.1. Constraints from the diffuse Galactic ISM used by Zubko et al. (2004): (left) Abundances (centre) UV-visible extinction curve (right) IR emission. In each plot, the "model" curve corresponds to what is obtained for the final size distribution. In the (centre) and (right) panels the plain, dashed and dotted lines are the PAH, graphite and silicate contributions to the model curve.

Table 7.1. Properties of the dust components (adapted for our dust composition from Table 2 of Zubko et al. 2004).

| Constituent | Composition | Mass density $\left[\mathrm{g} . \mathrm{cm}^{-3}\right]$ | $a_{\min }[\mu \mathrm{m}]$ | $a_{\max }[\mu \mathrm{m}]$ | References for Optical properties |
| :--- | :---: | :---: | :---: | :---: | :---: |
| PAHs | C | 2.24 | $3.5 \times 10^{-4}$ | $5 \times 10^{-3}$ | Li \& Draine $(2001)^{a}$, Draine \& Li $(2007)^{b}$ |
| Graphite | C | 2.24 | $3.5 \times 10^{-4}$ | 0.33 | Laor \& Draine (1993) |
| Silicate | $\mathrm{MgFeSiO}_{4}$ | 3.5 | $3.5 \times 10^{-4}$ | 0.37 | Weingartner \& Draine (2001) |

[^36]
## Optical properties

Now that we have a size distribution, we need to compute the absorption efficiencies, $Q_{a b s}^{i}(a, \lambda)$, of each of our dust species. These optical properties are derived using a Mie code and the method presented in Laor \& Draine (1993). The PAH optical properties have been updated by Draine \& Li (2007) who use more accurate band profiles constrained by Spitzer spectra compared to Li \& Draine (2001). In the PAHs component we distinguish between the neutral and ionised PAHs. The absorption efficiencies are shown in Fig. 7.2 for the four dust species: neutral PAHs, ionised


Fig. 7.2. Grain properties: (top) Size distributions from Zubko et al. (2004) for PAHs (red), graphite (green) and silicate grains (blue) (bottom) Absorption efficiencies, normalised to a given grain radius $a, Q_{a b s}(\lambda) / a$, for neutral PAHs (top left panel), ionised PAHs (top right panel), graphite (bottom left panel) and silicate (bottom right panel) grains. The colours indicate different values for the grain radius $a$.

PAHs, graphite and carbon grains for different grain radius $a$. The opacity of the adopted dust composition, $\kappa_{a b s}$, is the sum of the opacity for each dust specie: PAHs, carbon and silicate grains. Following Eq. 1.13 the opacity for the $\mathrm{i}^{\text {th }}$ dust specie is given by:

$$
\begin{equation*}
\kappa_{a b s}^{i}(\lambda)=\frac{\int \pi a^{2} Q_{a b s}^{i}(a, \lambda) G_{i}(a) \mathrm{d} a}{\int \frac{4}{3} \pi \rho_{i} a^{3} G_{i}(a) \mathrm{d} a} \tag{7.9}
\end{equation*}
$$

The submm emissivity index $\beta$ is defined by the logarithmic submm slope of the absorption efficiency (see Section 1.2.3) or equivalently of the opacity:

$$
\begin{equation*}
Q_{a b s}(\lambda)=Q_{a b s}\left(\lambda_{0}\right)\left(\frac{\lambda_{0}}{\lambda}\right)^{\beta} \leftrightarrow \kappa_{a b s}(\lambda)=\kappa_{a b s}\left(\lambda_{0}\right)\left(\frac{\lambda}{\lambda_{0}}\right)^{-\beta} \tag{7.10}
\end{equation*}
$$

Fig. 7.3 shows the total opacity for our dust mixture. In our case, the fit of the submm part of the opacity gives an emissivity index $\beta=2.0$ (Fig. 7.3), where we chose to normalise at $\lambda_{0}=160 \mu \mathrm{~m}$ with $\kappa_{a b s}\left(\lambda_{0}\right)=1.4 \mathrm{~m}^{2} \mathrm{~kg}^{-1}$.


Fig. 7.3. Opacities for the adopted dust composition (in black) and for each dust specie: silicates in blue, graphite in green, neutral PAHs in red and ionised PAHs in orange. The $\kappa_{a b s}$ for each specie has been calculated using Eq. 7.9 and $f_{P A H+}^{n}, f_{P A H 0}^{n}, f_{g r}^{n}$ and $f_{s i l}^{m}$ are the corresponding mass fractions (see Section 7.1.3). The total opacity for our dust mixture (in black) is the sum of the different contributions displayed in colours. The $\lambda^{-2}$ dependence of the submm part of the total opacity is shown in grey.

Note that when we say "grain properties" we encompass two things: the size distribution and the optical properties. The optical properties are intrinsic properties of the grains. The size distribution, however, assumes certain abundances for the elements constituting the dust and a given
illumination (solar abundances and Galactic ISRF in our case). Technically this size distribution is not directly applicable to other galaxies where abundances and illumination conditions differ from the diffuse Galactic ISM (especially for dwarf galaxies, see Section 2.2.2). Ideally we would need to reproduce the method presented by Zubko et al. (2004) for each galaxy we want to model and get a size distribution that simultaneously reproduces the elemental abundances, UV extinction, and IR emission of the considered galaxy. However, the lack of observational constraints does not enable us to follow such method, and, in order to move forward we approximate the grain size distribution in the galaxy we want to model by that from the BARE-GR-S model of Zubko et al. (2004).

### 7.1.3 One ISM mass element

The first step of the model is to compute the SED emitted by a single mass-element of the ISM, $l_{\nu}^{\text {dust }}(U, \lambda)$, illuminated by a stellar radiation field with uniform intensity $U . l_{\nu}^{\text {dust }}(U, \lambda)$ is given by:

$$
\begin{equation*}
l_{\nu}^{\text {dust }}(U, \lambda)=f_{P A H}^{m}\left[f_{P A H+}^{m} l_{\nu}^{P A H+}(U, \lambda)+\left(1-f_{P A H+}^{m}\right) l_{\nu}^{P A H 0}(U, \lambda)\right]+f_{g r}^{m} l_{\nu}^{g r}(U, \lambda)+f_{s i l}^{m} l_{\nu}^{s i l}(U, \lambda) \tag{7.11}
\end{equation*}
$$

where:

- $f_{P A H+}^{n}$ is the mass fraction of ionised PAHs compared to the total PAHs mass,
- $f_{P A H}^{n}$ the mass fraction of total PAHs (neutral and ionised),
- $f_{g r}^{n}$ the mass fraction of graphite grains, and
- $f_{s i l}^{n}$ is the mass fraction of silicate grains.

We have $f_{P A H}^{n}+f_{g r}^{n}+f_{s i l}^{n}=1$. $l_{\nu}^{P A H+}(U, \lambda), l_{\nu}^{P A H 0}(U, \lambda), l_{\nu}^{g r}(U, \lambda)$ and $l_{\nu}^{s i l}(U, \lambda)$ are the corresponding size distribution integrated monochromatic luminosities per unit dust mass.

Following Eq. 1.37, the monochromatic luminosities per unit dust mass are given $\mathrm{by}^{5}$ :

$$
\begin{equation*}
l_{\lambda}^{i}(\lambda)=\frac{1}{M_{i}} \int 4 \pi^{2} a^{2} Q_{a b s}^{i}(\lambda, a) S_{\lambda}^{i}(\lambda, a) G_{i}(a) \mathrm{d} a \tag{7.12}
\end{equation*}
$$

where $M_{i}$ is the mass of the $\mathrm{i}^{\text {th }}$ dust specie and is given by: $M_{i}=\int \frac{4}{3} \pi a^{3} \rho_{i} G_{i}(a) \mathrm{d} a$.
Note that this time in Eq. 7.12, compared to Eq. 7.3, $G_{i}(a)$ is known and we want to compute the $l_{\lambda}^{i}$ using the optical properties computed in 7.1.2.

However, as mentioned in Section 7.1.2, the assumed size distribution $G_{i}(a)$ is the size distribution of dust grains in the Galaxy. In another galaxy, the grain size distribution could be different. Schematically, a larger number of SN shocks or very intense radiation field will tend to erode the bigger grains, transferring mass into smaller grains. On the contrary, in dense and cold environments, grains can accrete and coagulate thus transferring mass from small to big grains. Thus we separate the graphite and silicate grains into very small grains (VSGs) and big grains (BGs), to account for an eventual variation of the proportion of small grains compared to the proportion of big grains. The transition size between the two is fixed at $a_{t}=0.01 \mu \mathrm{~m}$. A larger number of small grains compared to that in the Galaxy will appear as a higher NIR continuum in the modelled SED. The last two members of the right hand side of Eq. 7.11 thus become:

$$
\begin{equation*}
f_{g r}^{m} l_{\nu}^{g r}(U, \lambda)=f_{V S G, g r}^{m} l_{\nu}^{V S G, g r}(U, \lambda)+f_{B G, g r}^{m} r_{\nu}^{B G, g r}(U, \lambda) \tag{7.13}
\end{equation*}
$$

[^37]and:
\[

$$
\begin{equation*}
f_{s i l}^{m} l_{\nu}^{s i l}(U, \lambda)=f_{V S G, s i l}^{m} l_{\nu}^{V S G, s i l}(U, \lambda)+f_{B G, s i l}^{m} l_{\nu}^{B G, s i l}(U, \lambda) \tag{7.14}
\end{equation*}
$$

\]

with: $f_{P A H}^{n}+f_{V S G, g r}^{n}+f_{B G, g r}^{n}+f_{V S G, s i l}^{n}+f_{B G, s i l}^{n}=1$. The integration on the size distribution in Eq. 7.12 is performed over the $\left[a_{i, \text { min }}, a_{t}\right]$ range to get $l_{\nu}^{V S G, i}(U, \lambda)$; and over the $\left[a_{t}, a_{i, \text { max }}\right]$ range to get $l_{\nu}^{B G, i}(U, \lambda)$ with $\mathrm{i}=\{\mathrm{gr}$, sil $\}$.

The transition radius $a_{t}$ implies that the VSGs always undergo stochastic heating, and that BG are either stochastically heated or in thermal equilibrium. The source functions $S_{\lambda}^{i}(\lambda, a)$ are computed accordingly:

- For BGs at thermal equilibrium, $S_{\lambda}^{i}(\lambda, a) \simeq \mathrm{B}_{\lambda}\left(\lambda, T_{d}(a)\right)$, where $T_{d}(a)$ is determined from Eq. 7.5.
- For stochastically heated VSGs or BGs, their temperature probability distributions are computed following the method of Guhathakurta \& Draine (1989).

Fig. 7.4 shows the temperature probability distributions for the different grain species for different grain radiuses $a$, and the enthalpies ${ }^{6}$ used in their derivation. We see that large graphite and silicate grains have a narrow temperature probability distribution: their temperature does not oscillate much around their equilibrium temperatures. PAHs, small graphite grains and small silicate grains have, on the contrary, broad temperature distributions, showing the large range of temperatures these grains can span.

Here we assume that the ISRF heating the grains has the spectral shape of the solar neighbourhood (Mathis et al. 1983), and we parameterise its mean intensity, $J_{\nu}$, by:

$$
\begin{equation*}
J_{\nu}(\lambda)=U \times J_{\nu}^{\text {Gal }}(\lambda) \tag{7.15}
\end{equation*}
$$

where $J_{\nu}^{G a l}(\lambda)$ is the mean intensity of the ISRF from Mathis et al. (1983), normalised to the solar neighbourhood value: $4 \pi \times \int J_{\nu}^{G a l} \mathrm{~d} \nu=2.2 \times 10^{5} \mathrm{~W} . \mathrm{m}^{-2}$. When $\mathrm{U}=1$, it means that the ISM mass element is illuminated by a stellar radiation field having the intensity of the solar neighbourhood ISRF from Mathis et al. (1983).

We now have all of the elements to generate dust SED templates for a mass-element of the ISM with this dust composition at a given intensity of the ISRF. We generate these templates for a grid of starlight intensities, $U$, from $10^{-2}$ to $10^{7}$.

### 7.1.4 Combining several ISM mass elements

Now that we have the emission from a single ISM mass-element, we want to find the proper combination of mass-elements, i.e., of illumination conditions, that fits the observations we have for our sources. Indeed the emission from a galaxy will likely be a combination of different regions with different physical conditions. Here, as mentioned in 7.1.1, we assume that the dust properties are uniform throughout the galaxy, i.e., the size distribution and the optical properties are constant within the galaxy, and that only the starlight intensity varies.

The total specific monochromatic luminosity emitted by dust in the galaxy is given by:

$$
\begin{equation*}
L_{\nu}^{\text {tot }, \text { dust }}=\int l_{\nu}^{\text {dust }}(U(M), \lambda) \mathrm{d} M \tag{7.16}
\end{equation*}
$$

or:

[^38]

Fig. 7.4. (top): Enthalpies for the considered dust species. (bottom row) Temperature probability distributions for the different grain species. The colour codes the grain size: from black (small radius) to red (big radius).

$$
\begin{equation*}
L_{\nu}^{t o t, d u s t}=\int\left(\sum_{i=1}^{4} f_{i}^{m} l_{\nu}^{i}\right) \frac{d M}{d U} \mathrm{~d} U \tag{7.17}
\end{equation*}
$$

where $\mathrm{i}=\{\mathrm{PAH}+$, PAH0, gr, sil $\}$ (see Eq. 7.11).
We use the empirical prescription from Dale et al. (2001a) to describe the distribution of starlight intensities per unit dust mass: it can be approximated by a power law of index $\alpha$ :

$$
\begin{equation*}
\frac{d M}{d U}=c s t \times U^{-\alpha} \tag{7.18}
\end{equation*}
$$

In most cases, this prescription is flexible enough to reproduce dense and diffuse media, and provides a simple parameterization of the physical conditions in the ISM. With the total dust mass, given by:

$$
\begin{equation*}
M_{d u s t}=\int_{U_{\min }}^{U_{\min }+\Delta U} \frac{d M}{d U} \mathrm{~d} U \tag{7.19}
\end{equation*}
$$

we get:

$$
\begin{equation*}
M_{\text {dust }}=\frac{c s t}{1-\alpha}\left(U_{\max }^{1-\alpha}-U_{\min }^{1-\alpha}\right) \tag{7.20}
\end{equation*}
$$

with $U_{\max }=U_{\min }+\Delta U$. Let $\mathrm{K}=\frac{U_{\max }^{1-\alpha}-U_{\min }^{1-\alpha}}{1-\alpha}$, we have $c s t=M_{\text {dust }} / K$. Injecting this into Eq. 7.17 gives:

$$
\begin{equation*}
L_{\nu}^{t o t, \text { dust }}=\frac{M_{\text {dust }}}{K} \int\left(\sum_{i=1}^{3} f_{i} l_{\nu}^{i}\right) U^{-\alpha} \mathrm{d} U \tag{7.21}
\end{equation*}
$$

Emission from old stars can also contribute to the IR emission and especially in the NIR. Thus we add a stellar continuum to the dust emission, parameterised by the stellar mass of the galaxy $M_{\star}$ :

$$
\begin{equation*}
L_{\nu}^{\star}=M_{\star} l_{\nu}^{\star} \tag{7.22}
\end{equation*}
$$

$l_{\nu}^{\star}$ is the specific monochromatic luminosity per unit stellar mass of a 1 Gyr stellar population, synthesised with the PEGASE ${ }^{7}$ model (Fioc \& Rocca-Volmerange 1997), assuming a Salpeter Initial Mass Function (IMF): $N(M) d M=M^{-1.35} d M$ and an initial metallicity $\mathrm{Z}=Z_{\odot}$. The age of the stellar population is not of great importance here as this population will essentially be constrained by the NIR bands up to IRAC $4.5 \mu \mathrm{~m}$. However, the stellar mass is poorly determined and does not give a reliable estimate of the true stellar mass of the galaxy. The point of this stellar component is to have a good fit in the NIR, in order to extrapolate the stellar contribution in the MIR.

The total emission from the galaxy is thus:

$$
\begin{equation*}
L_{\nu}^{\text {tot }}=L_{\nu}^{\text {dust }, \text { tot }}+L_{\nu}^{\star} \tag{7.23}
\end{equation*}
$$

For each band at a wavelength $\lambda_{i}$ we can compute the synthetic luminosity, $L_{\nu}^{\text {model }}\left(\lambda_{i}\right)$ by convolving the model with the RSRF of each band and using the appropriate spectral convention of each instrument. The various $L_{\nu}^{\text {model }}\left(\lambda_{i}\right)$ are computed using Eqs. 6.8, 6.16, 6.18 and 6.19.

The fitting procedure is an iterative process where we combine all of the templates of one ISM mass-element we need to fit the observations and stops when observational constraints are

[^39]minimized. The fit is performed with the mpcurvefit procedure in IDL based on the LevenbergMarquardt methods.

### 7.1.5 Limitations of the model

This model has been built on a certain number of assumptions, and thus presents some limitations. We argue here that these limitations do not biais our final results.

Geometry : Ideally, radiative transfer models are a more rigorous way of reconstructing the total emission of a galaxy. In these models, the emission from a mass-element of the ISM is computed via a radiation transfer model, and then is propagated to the observer. In order to do so, we need to know about the geometry of the stellar sources and of the dust and gas within the galaxy. In our case, we do not have access to such level of detail for all of our galaxies. Moreover we are interested by total dust masses and the dust mass is dominated by the mass of the large dust grains, in equilibrium with the radiation field of the galaxy. The spectrum of these large dust grains is determined by the stellar power absorbed by the grains and not by the spatial distribution of the stellar sources.

Local variations of grain properties : Dust grains emissivity values could vary in different regions of the modelled galaxy. In this case, our model would lead to a misestimate of the dust mass in regions where the emissivity differs from the emissivity assumed here. However these variations are very poorly known, and difficult to constrain. Grain-grain coagulation for example could be responsible for a local increase of the FIR opacity in dense clouds (Stepnik et al. 2003). The emissivity of the dust grain could also increase with the grain temperature (Meny et al. 2007), and thus we would overestimate the dust mass in hot dust regions. However here we are interested in dust masses for total galaxies, and not in local variations of the dust mass within these galaxies. Thus we believe that the effects described here would not have too much impact on the total dust mass (i.e., if they have any, it would affect the dust mass within its error bar).

Radiation field : In this model, we assume that the ISRF has the shape of the Galactic diffuse ISM. However, some studies (e.g., Madden et al. 2006) have shown that the ISRF of dwarf galaxies differs from the radiation field of "normal" more metal-rich galaxies. The ISRF in low-metallicity dwarf galaxies is often more intense and harder, due to the predominance of massive stars, and a larger mean free path length of the ionizing photons, attributed to the lower dust attenuation in these galaxies. The shape of the radiation field determines the emission of out-of-equilibrium small grains. Increasing the hardness of the radiation field increases the maximum temperature the small grains can reach when they fluctuate. However, these small grains only have a minor contribution to the total dust mass, and thus the assumed shape of the ISRF does not bias our estimation of the total dust mass.

### 7.2 Modelling the DGS and KINGFISH galaxies

The model described in the previous Section is applied to the DGS and KINGFISH galaxies. For the DGS galaxies, the observational constraints are from 2MASS, Spitzer, WISE, IRAS (presented in Chapter 6) and Herschel instruments (presented in Chapter 4). For the KINGFISH galaxies, the observational constraints originate from the same instruments (except WISE and Spitzer/IRS).

Some galaxies are not detected at one or several wavelengths. We do not consider the upper limits in the fits by imposing a zero weight on the points and imposing to the best fit model to be consistent with the non-detections.

In order to study the submm excess in more details in Section 8.3, we do not include in the fit the SPIRE $500 \mu \mathrm{~m}$ point nor any available ground-based submm data (i.e., we put a zero weight on these points). Indeed, the submm excess can already be observed for some dwarfs galaxies at this wavelength, thus trying to fit this point can, in some cases, bias our estimation of the dust mass in these galaxies.

### 7.2.1 Setting the parameters

In our case, the model can be described by 8 parameters:

- $M_{\text {dust }}$ : the total dust mass,
- $U_{m i n}$ : the minimum of the starlight intensity distribution,
- $\Delta U$ : the difference between the maximum and minimum of the starlight intensity distribution,
- $\alpha$ : the index of the power law describing the starlight intensity distribution,
- $f_{P A H}$ : the PAHs mass fraction, normalised to the Galactic PAH mass fraction (when $f_{P A H}=0.5$ it means that the mass fraction of PAH in the modelled galaxy is half the mass fraction of PAHs in the Galaxy),
- $f_{i o n}$ : the ionised-to-total PAH mass ratio (when $f_{i o n}=0.5$ it means that $50 \%$ of the PAHs are ionised),
- $f_{v s g}$ : the very small grains mass fraction, i.e., the mass fraction of grains with radius $a \leq 10$ nm , normalised to the Galactic value (when $f_{v s g}=0.5$ it means that the mass fraction of very small grains in the modelled galaxy is half the mass fraction of very small grains in the Galaxy), and
- $f_{s i l}$ : the silicate-to-(silicate+graphite) grains mass fraction (when $f_{s i l}=0.5$ it means that there is the same mass of silicate and graphite grains in the ISM.)

The different mass fractions $f_{P A H}^{n}, f_{P A H+}^{n}, f_{V S G, g r}^{n}, f_{B G, g r}^{n}, f_{V S G, s i l}^{n}$ and $f_{B G, s i l}^{m}$ can be computed from the above parameters and Galactic mass fractions:

$$
\left\{\begin{align*}
f_{P A H}^{m} & =f_{P A H, \odot}^{m} \times f_{P A H}  \tag{7.24}\\
f_{P A H+}^{m} & =f_{P A H, \odot}^{m} \times f_{P A H} \times f_{i o n} \\
f_{V S G, g r}^{m} & =\left(1-f_{P A H}^{m}\right) \times\left(1-f_{\text {sil }}\right) \times f_{V S G, g r, \odot} f_{v s g} \\
f_{B G, g r}^{m} & =\left(1-f_{P A H}^{m}\right) \times\left(1-f_{s i l}\right) \times\left(1-f_{V S G, g r, \odot} f_{v s g}\right) \\
f_{V S G, s i l}^{m} & =\left(1-f_{P A H}^{m}\right) \times f_{s i l} \times f_{V S G, s i l, \odot} f_{v s g} \\
f_{B G, s i l}^{m} & =\left(1-f_{P A H}^{m}\right) \times f_{s i l} \times\left(1-f_{V S G, s i l, \odot} f_{v s g}\right)
\end{align*}\right.
$$

where $f_{P A H, \odot}^{m}=0.0457$ is the Galactic PAH mass fraction; $f_{V S G, g r, \odot}=0.1846$ is the proportion of graphite VSGs in the Galaxy; and $f_{V S G, s i l, \odot}=0.1655$ is the proportion of silicate VSGs in the Galaxy. These values can be obtained from the dust properties given by Zubko et al. (2004).

We fix the $f_{\text {sil }}$ parameter to its Galactic value, leaving us with 7 free parameters:

$$
\begin{equation*}
f_{s i l}=\frac{f_{s i l, \odot}^{m}}{f_{s i l, \odot}^{m}+f_{g r, \odot}^{m}}=\frac{0.6596}{0.6596+0.2947}=0.6912 \tag{7.25}
\end{equation*}
$$

where $f_{s i l, \odot}^{m}$ and $f_{g r, \odot}^{m}$ are the Galactic mass fraction of silicate and graphite grains respectively, taken from Zubko et al. (2004).

As IRS spectra are available for the DGS sample, we can put constraints on the $f_{i o n}$ and $f_{P A H}$ parameters. When no IRS constraints are available, we leave these parameters free. However, whenever PAH features are absent from the IRS spectrum, we fix $f_{P A H}=0$. When there are no PAHs, the value of $f_{i o n}$ does not have any importance and so we fix it arbitrarily to $f_{i o n}=0.5$. If the IRS spectrum provides good SL constraints on the ionised PAH lines we leave this parameter free, otherwise, we also fix $f_{\text {ion }}=0.5$. These choices for $f_{\text {ion }}$ and $f_{P A H}$ are done in order to improve the quality of the fit. The value of $f_{i o n}$ has no influence on the dust mass as it controls the proportion of ionised PAHs.

For some DGS galaxies, the MIR continuum shape outlined by the IRS spectrum cannot be well fitted by our model. In these cases, we add an extra modified blackbody component in the MIR, with a fixed $\beta=2.0$ and a temperature varying between 80 and 300 K . It can be physically interpreted as the contribution from a hot Hir region to the total emission of the galaxy. As dwarf galaxies have small physical sizes and low dust attenuation, the emission from the energetic Hir regions can indeed affect the total emission from the whole galaxy. In these cases, this tells us that the Dale et al. (2001a) prescription of $\frac{d M}{d U} \propto U^{-\alpha}$ may not be flexible enough to provide a satisfactory fit for these galaxies. This can have a small impact on the dust mass, as we are adding warm equilibrium dust grains (Fig. 7.5). However, we demonstrate in Section 7.2.4 for cases where the impact on the dust mass is important (i.e., greater than the dust mass error bar, red points in Fig. 7.5) that the addition of a warm modified blackbody is necessary to have a coherent MIR to submm shape of the SED.


Fig. 7.5. Comparison of the dust masses estimated with an additional warm MIR blackbody $\left(M_{d u s t}(w B B)\right)$ compared to the dust masses obtained without this extra-component $\left(M_{d u s t}(w o B B)\right)$ for galaxies for which this additional blackbody is needed (see Section 7.2 .1 ). The colours show whether $M_{\text {dust }}(w B B)$ and $M_{\text {dust }}(w o B B)$ are consistent within errors (in green) or not (in red).

### 7.2.2 Extracting total infrared luminosities, mean radiation field intensities and star-formation rates

From the best fit model we can extract two quantities derived from the parameters for the galaxy: the TIR luminosity $L_{T I R}$ and the mass-averaged starlight intensity $\langle U\rangle$. These two parameters are given by:

$$
\begin{equation*}
L_{T I R}=\int_{0}^{\infty} L_{\nu}^{\text {tot }, d u s t}(\nu) \mathrm{d} \nu=\int_{1 \mu m}^{1000 \mu m} L_{\nu}^{\text {tot }, d u s t}(\nu) \mathrm{d} \nu \tag{7.26}
\end{equation*}
$$

and:

$$
\begin{equation*}
<U>=\frac{1}{M_{\text {dust }}} \int_{U_{\min }}^{U_{\max }} U \times \frac{d M_{\text {dust }}}{d U} \mathrm{~d} U \tag{7.27}
\end{equation*}
$$

Eq. 7.27 translates into:

$$
<U>= \begin{cases}\frac{1-\alpha}{2-\alpha} \frac{U_{\max }^{2-\alpha}-U_{\min }^{2-\alpha}}{U_{\max }^{2-\alpha}-U_{\min }^{2}} & \text { if } \alpha \neq 1 \& \alpha \neq 2  \tag{7.28}\\ \ln \left(1+\Delta U U U_{\min n}\right. & \text { if } \alpha=1 \\ U_{\min } \ln \left(1+\Delta U / U_{\text {min }}\right)\left(\frac{U_{\min +\Delta U}}{\Delta U}\right) & \text { if } \alpha=2\end{cases}
$$

From the $L_{T I R}$ we can also compute an estimation of the star-formation rate (SFR) for the galaxies in our sample, using the formula of Kennicutt (1998) presented in Eq. 3.22.

### 7.2.3 Errors

As in Chapter 5 the errors on the various parameters are estimated by generating 300 random realisations of the SEDs within errors in order to get a distribution for the parameters. For each galaxy, we randomly perturb our fluxes within the errors bars and perform fits of the perturbed SEDs. Here again we must take special care for errors which are correlated between different bands. We detail here the decomposition of the errors for all the instruments we use, except Herschel as this decomposition has already been presented in Section 5.2.2.

2MASS: Jarrett et al. (2003) quote a $2-3 \%$ uncertainty on the zero-magnitude flux values. To be conservative, we assume a non-correlated error of $3 \%$.

IRAC: The total calibration uncertainty used for the DGS and KINGFISH IRAC fluxes is $\sim 10 \%$. This can be decomposed into two parts:

- Reach et al. (2005) give a $2 \%$ uncertainty in all of the IRAC bands. This error is correlated between the four bands.
- The IRAC Instrument Handbook (Section 4.3) also recommends using a $10 \%$ error to account for several systematic effects in the calibration. This error is independent from band to band.

IRS: The IRS spectra extracted from the CASSIS database already provide a decomposition of the total error on the flux densities into three parts:

- Part of the error is the peak-to-peak uncertainty and represents the measurement error, independent for the different wavelengths.
- Part of the error is due to the flux difference between the two nod spectra and is correlated for all of the wavelengths over both SL and LL ranges.
- The third component of the error in IRS spectra is the correlation error. It is not clear whether this error is correlated between the two SL and LL modules, so, to be conservative, we assume that the correlation errors are correlated for all of the SL wavelengths on one side and for all of the LL wavelengths on the other side. According to Lebouteiller et al. (2011), the global IRS calibration is better than the $2 \%$ level.

MIPS: For the DGS MIPS photometry, Bendo et al. (2012) used a $4 \%$ calibration error at $24 \mu \mathrm{~m}$ (Engelbracht et al. 2007), $10 \%$ at $70 \mu \mathrm{~m}$ (Gordon et al. 2007) and $12 \%$ at $160 \mu \mathrm{~m}$ (Stansberry et al. 2007). For the KINGFISH, the same uncertainties were used by Dale et al. (2007) except for MIPS $70 \mu \mathrm{~m}$ where they adopted a $7 \%$ calibration error. According to the MIPS Instrument Handbook, the calibration of the MIPS $160 \mu \mathrm{~m}$ band has been done using the $24 \mu \mathrm{~m}$ and $70 \mu \mathrm{~m}$ observations of asteroids. We can thus consider that the calibration errors for MIPS $24 \mu \mathrm{~m}$ and MIPS $70 \mu \mathrm{~m}$ are independent but that they are correlated with MIPS $160 \mu \mathrm{~m}$.

WISE: WISE calibration has been performed on stars and is tied to Spitzer calibration according to Jarrett et al. (2011). For each wavelength, the correlations between the bands can be summarised this way:

- The WISE $3.4 \mu \mathrm{~m}$ calibration error is decomposed in an independent part, proper to WISE, of $2.4 \%$, and is correlated with the IRAC $3.6 \mu \mathrm{~m}$ band. Thus we add the IRAC $3.6 \mu \mathrm{~m}$ calibration uncertainty to the WISE $3.4 \mu \mathrm{~m}$ calibration error.
- Following the WISE $3.4 \mu \mathrm{~m}$ band, the WISE $4.6 \mu \mathrm{~m}$ error has an independent part of $2.8 \%$ and is correlated with the IRAC $4.5 \mu \mathrm{~m}$ band.
- Similarly, the WISE $12 \mu \mathrm{~m}$ error has an independent part of $4.5 \%$ and is correlated with the IRS-SL/LL modules.
- And finally, the WISE $22 \mu \mathrm{~m}$ error has an independent part of $5.7 \%$ and is correlated with the MIPS $24 \mu \mathrm{~m}$ band.

IRAS: According to the IRAS explanatory supplement ${ }^{8}$, the calibration of $\operatorname{IRAS}$ has been tied to the Rieke et al. (1984) ground-based photometric system at $12 \mu \mathrm{~m}$. The three IRAS bands at 12,25 and $60 \mu \mathrm{~m}$ have been calibrated using stellar models, and the IRAS $100 \mu \mathrm{~m}$ calibration used asteroids. The relative uncertainties relative to the ground-based $12 \mu \mathrm{~m}$ are $2 \%, 5 \%$ and $5 \%$ for IRAS 12, 25 and $60 \mu \mathrm{~m}$ respectively, independent from band to band. The absolute uncertainty on the $12 \mu \mathrm{~m}$ flux density is $4 \%$, in common and correlated between the three bands. The uncertainty at $100 \mu \mathrm{~m}$ is $10 \%$, and is not correlated with any of the other IRAS bands.

SCUBA for KINGFISH: Dale et al. (2005) give a calibration uncertainty of $25 \%$ and $15 \%$ for the 450 and $850 \mu \mathrm{~m}$ bands respectively. We assume that these errors are independent between the two bands.

[^40]Other ground-based instruments: For the other submm measurements (SCUBA, LABOCA and MAMBO), the calibration is derived directly from the observations as it depends on the conditions during the observations. It will be different from galaxy to galaxy. The values for the calibration errors can be found in the references quoted for the submm flux densities in Table 6.12. The errors from the different instruments are independent from one another.

As in Section 5.2.2, the perturbation of the observed fluxes is the sum of two components:

- A normal random independent variable representing the measurement errors at each wavelength.
- A normal random variable describing the calibration errors that takes into account the correlation between the wavebands as described above.

Here again we quote the $66.67 \%$ confidence level for our parameters.

### 7.2.4 Notes on individual DGS galaxies

- Haro 11 - As this galaxy has silicate absorption features around $10 \mu \mathrm{~m}$, we add extinction in our modelling. The output luminosity is modified by:

$$
\begin{equation*}
L_{\nu}^{\text {tot }, d u s t}\left(\lambda, A_{V}\right)=L_{\nu}^{\text {tot }, \text { dust }}\left(\lambda, A_{V}=0\right) \times e^{-\tau_{0}(\lambda) A_{V} / 1.086} \tag{7.29}
\end{equation*}
$$

where $\tau_{0}$ is the optical depth, determined from the optical properties and the size distribution of our dust composition:

$$
\begin{equation*}
\tau_{0}(\lambda)=\sum_{i=1}^{3} \frac{\int \pi a^{2} Q_{e x t}^{i}(a, \lambda) F_{i}(a) \mathrm{d} a}{\int \frac{4}{3} \pi \rho_{i} a^{3} F_{i}(a) \mathrm{d} a} \tag{7.30}
\end{equation*}
$$

The extinction parameter $A_{V}$ is left varying between 1 and 20 , and the best fit value is $A_{V}=5$. The fit is clearly improved in the NIR-MIR range after introducing $A_{V}$ (see Fig. 7.6). The stellar continuum is also affected by extinction and the PEGASE stellar template can not provide a reliable fit of the 2MASS NIR points. Instead, we replace it by a blackbody (fixing $\beta=0$ ) and imposing a temperature $T \geq 1000 \mathrm{~K}$. We obtain a best fit temperature $T=30000 \mathrm{~K}$, corresponding to a peak of the blackbody emission around $0.1 \mu \mathrm{~m}$, in the UV - a sign of quite a young stellar population. This is consistent with what is found by Adamo et al. (2010) who found a peak of cluster formation at $\sim 3.5 \mathrm{Myr}$ ago, and with the UV part of the SED presented in Cormier et al. (2012). Fig. 7.6 shows the improvement of this additional component compared to the default model.

- HS 0822+3542 - The PACS upper limits at 70 and $100 \mu \mathrm{~m}$ are not consistent at all with the rest of the MIR to FIR photometry (see Fig. 7.23). Thus we do not impose that these upper limits are respected in the fit. We decide to change the reported upper limits for the 70 and $100 \mu \mathrm{~m}$ PACS wavelengths and we recommend using the synthetic photometry provided by the model, i.e., $F_{70} \leq 41 \mathrm{mJy}$ and $F_{100} \leq 48 \mathrm{mJy}$. These values are reported in Table 4.1.
- HS 1236+3937- Only Spitzer IRAC and IRS, and Herschel PACS and SPIRE data are available for this galaxy, and the galaxy is not detected with any of these instruments. So we can not report a SED for this galaxy.
- HS 1304+3529 - To obtain a better fit at MIR wavelengths, we add an extra MIR modified blackbody ( $\beta$ fixed to 2.0 and best fit temperature $T=100 \mathrm{~K}$ ). Additionally, this extra MIR modified


Fig. 7.6. SED for Haro11 when no extinction is assumed (left) and with extinction (right). The colours and symbols for all of the SEDs presented in this Section are described in Fig.7.23.
blackbody improves the fit in the FIR. This is shown in Fig. 7.7.


Fig. 7.7. SED for HS1304+3529 without (left) and with (right) the extra MIR modified blackbody.

- HS 1442+4250 - There are only 4 detected points for this galaxy: IRAC $5.8 \mu \mathrm{~m}$, IRAC $8.0 \mu \mathrm{~m}$, MIPS $24 \mu \mathrm{~m}$ and PACS $70 \mu \mathrm{~m}$, and thus not enough constraints to fit a SED.
- HS 2352+2733 - There are only 3 detected points for this galaxy: MIPS $24 \mu \mathrm{~m}, \mathrm{PACS} 70 \mu \mathrm{~m}$ and PACS $100 \mu \mathrm{~m}$, and thus not enough constraints to fit a SED. However, we can get a rough estimate of the dust mass by fitting a modified blackbody with a fixed $\beta=2.0$, through these 3 points. We obtain a temperature of $\mathrm{T}=52 \mathrm{~K}$ and $\mathrm{M}_{\text {dust, } B B}=1.01 \times 10^{4} M_{\odot}$ (see Fig. 7.8). This mass can bee seen as a lower limit of the real dust mass (see Chapter 8). Taking the median ratio of the dust masses estimated with the full dust model to the dust masses estimated with a $\beta=2.0$ modified blackbody for the DGS ( 0.51 , see Section 8.2 .1 where this ratio is derived), we can estimate a total dust mass of $\sim \mathrm{M}_{\text {dust }}=1.98 \times 10^{4} M_{\odot}$.
- I Zw 18 - To get a better match of the IRS spectrum, we add an extra MIR modified blackbody ( $\beta$ fixed to 2.0 and best fit temperature $T=100 \mathrm{~K}$ ). This is shown in Fig. 7.9.


Fig. 7.8. Modified blackbody fit for $\mathrm{HS} 2352+2733$ with a fixed $\beta=2.0$.


Fig. 7.9. SED for IZw18 without (left) and with (right) the extra MIR modified blackbody.


Fig. 7.10. SED for IIZw40 without (left) and with (right) the extra MIR modified blackbody. On the left panel, we zoom into the $3-150 \mu \mathrm{~m}$ range with the IRS spectrum overlaid on the figure in dark blue.

- II Zw 40 - This galaxy has a good IRS spectrum. However, when compared to other MIR measurements, it seems that the Spitzer IRS-LL spectrum is not consistent with the rest of the MIR - FIR photometry (see Fig. 7.10). We thus looked for the ISOCAM spectrum for this galaxy. ISOCAM covers a 5.6 to $16.3 \mu \mathrm{~m}$ range. We match the ISOCAM spectrum with the photometry by applying the same rescaling process as for the IRS SL spectrum, i.e., find a spline going through the IRAC $8 \mu \mathrm{~m}$ and WISE $12 \mu \mathrm{~m}$ bands. The comparison of the ISOCAM and the IRS spectra shows that ISOCAM is consistent with the SL part of the IRS spectrum. However the slope of the LL spectrum does not seem coherent with the slope outlined by the end of the ISOCAM spectrum. This means that for this galaxy, a wavelength dependent correction might also be needed for the IRS LL part of the spectrum. Unfortunately, the 3 constraints we have in the MIR, (WISE $22 \mu \mathrm{~m}$, MIPS $24 \mu \mathrm{~m}$ and $\operatorname{IRAS} 25 \mu \mathrm{~m}$ ) do not properly sample the LL spectrum and thus we cannot derive this wavelength dependent correction. Thus we use the ISOCAM spectrum in the modelling for this galaxy. To properly match the end of the ISOCAM spectrum, we add an extra modified blackbody ( $\beta$ fixed to 2.0 and best fit temperature $T=124 \mathrm{~K}$ ). This is shown in Fig. 7.10.
- Mrk 1450 - To get a better match of the IRS spectrum, we add an extra MIR modified blackbody ( $\beta$ fixed to 2.0 and best fit temperature $T=100 \mathrm{~K}$ ). This is shown in Fig. 7.11.


Fig. 7.11. SED for Mrk1450 without (left) and with (right) the extra MIR modified blackbody.
-Mrk 153 - This galaxy has a silicate emission feature in its IRS spectrum around $10 \mu \mathrm{~m}$. This galaxy is known to harbour an AGN and this emission feature could be originating from hot small silicate grains in the accretion disk around the AGN. Thus we leave the $f_{\text {sil }}$ parameter free for this galaxy. We get $f_{s i l}=0.9104$ (i.e., 1.3 times the Galactic value).

- Mrk 930 - To get a better match of the IRS spectrum, we add an extra MIR modified blackbody ( $\beta$ fixed to 2.0 and best fit temperature $T=100 \mathrm{~K}$ ). This is shown in Fig. 7.12.
- NGC1140 and NGC 1569 - The IRS SL slit only covers a small part of these galaxies. As they both are extended, we wonder if the SL rescaling is consistent with the rest of the NIR - MIR spectrum. We use the ISOCAM spectrum for comparison after applying the same rescaling step. The two spectra are consistent with each other and with the general NIR - MIR shape of the SED for both galaxies. Thus we consider that the IRS spectrum is reliable and we use it in the modelling. In the case of NGC1569, we also add an extra MIR modified blackbody ( $\beta$ fixed to 2.0 and best fit temperature $T=100 \mathrm{~K}$ ), to get a better match of the IRS spectrum. This is shown in Fig. 7.13.


Fig. 7.12. SED for Mrk930 without (left) and with (right) the extra MIR modified blackbody.


Fig. 7.13. Zoom on the SED between 3 and $150 \mu \mathrm{~m}$ for NGC1140 (left) where the ISOCAM spectrum is overlaid on the SED in light green. Same for NGC1569 (right). See Fig. 7.23 for the final SEDs.

- NGC 5253 - To get a better match of the IRS spectrum, we add an extra MIR modified blackbody ( $\beta$ fixed to 2.0 and best fit temperature $T=100 \mathrm{~K}$ ). This is shown in Fig. 7.14.
- NGC 6822 - If we leave the $f_{v s g}$ parameter free, this leads to unrealistically prominent PAH features for this galaxy. Thus we fix the $f_{v s g}$ parameter to the Galactic value for this galaxy. This leads to an increase of the dust mass of $\sim 15 \%$ which is within the $\sim 112 \%$ error bar ${ }^{9}$. This is shown in Fig. 7.15.
- Pox 186 - To get a better match of the IRS spectrum, we add an extra MIR modified blackbody ( $\beta$ fixed to 2.0 and best fit temperature $T=137 \mathrm{~K}$ ) (see Fig. 7.16). Pox 186 has a very broad FIR peak of the SED, and is not detected beyond $250 \mu \mathrm{~m}$. It has a very uncommon SED shape and a limited number of constraints. It looks like a very wide modified blackbody, i.e., a giant HıI region, that should be surrounded by some diffuse colder dust. Thus we introduce a single starlight intensity component, to mimic the emission of a colder diffuse ISM around this giant Hir region.

[^41]

Fig. 7.14. SED for NGC5253 without (left) and with (right) the extra MIR modified blackbody.


Fig. 7.15. SEDs for NGC6822 with $\mathrm{f}_{v s g}$ free (left) and fixed (right).


Fig. 7.16. SED for Pox186 with the extra diffuse component (in cyan), without (left) and with (right) the extra MIR modified blackbody.

We get a best fit $U=0.40$.

- SBS 0335-052 - This galaxy has a very surprising SED with a IR peak around 15-30 $\mu \mathrm{m}$, and an IRS spectrum showing silicate absorption features superimposed on a featureless continuum (see also Thuan et al. 1999b). The silicate absorption feature around $10 \mu \mathrm{~m}$ observed in the IRS spectrum indicates that extinction is prominent in this galaxy. Thus we add extinction in our model as for Haro 11, and get $A_{V}=1.1$. The peculiar shape of the IRS spectrum requires two MIR modified blackbodies to get a satisfactory fit ( $\beta$ fixed to 2.0 and best fit temperatures $T_{1}=281 \mathrm{~K}$ and $T_{2}=$ 125 K ). This is shown in Fig. 7.17.


Fig. 7.17. (left) SED for SBS0335-052 without extinction nor the additional MIR modified blackbodies. On the right panel, extinction and two extra MIR modified blackbodies have been added to improve the quality of the fit in the NIR - MIR range.

- SBS $1533+574$ - To get a better match of the IRS spectrum, we add an extra MIR modified blackbody ( $\beta$ fixed to 2.0 and best fit temperature $T=100 \mathrm{~K}$ ). This is shown in Fig. 7.18.


Fig. 7.18. SED for SBS1533+574 without (left) and with (right) the extra MIR modified blackbody.

- Tol 0618-402 - The longest wavelength detected for this galaxy is $12 \mu \mathrm{~m}$ in the WISE band. As we do not have any constraints in the FIR to get an accurate dust mass, we do not report a SED
for this galaxy.
- Tol 1214-277 - The MIPS $70 \mu \mathrm{~m}$ point is not coherent with the rest of the photometry: IRS and MIPS $24 \mu \mathrm{~m}$ on one side and PACS 70 and $100 \mu \mathrm{~m}$ on the other side (see Fig. 7.23). It is another confirmation that this point is discrepant (see Chapter 6, Section 6.2.3), and thus we do not consider it in the modelling. Similarly to Pox 186, Tol 1214-277 has a very broad SED peak in the FIR, resembling a giant Hir region. Here again we add a single starlight intensity component, and get a best fit $U=19.5$.
- UGC 4483 - The IRAC upper limit at $5.8 \mu \mathrm{~m}$ and the PACS upper limit at $160 \mu \mathrm{~m}$ do not seem consistent at all with the rest of the NIR to FIR photometry (see Fig. 7.23). Thus we do not impose that these upper limits are respected in the fit. We advise using the synthetic photometry provided by the model, i.e., $F_{5.8} \leq 0.85 \mathrm{mJy}$ and $F_{160} \leq 72 \mathrm{mJy}$. These flux densities are reported in Tables 4.1 and 6.9.
- UM 133 - This galaxy is not detected in the IRAC bands. However we still include the upper limits in the fit, in order to have (even approximative) constraints in the NIR (see Fig. 7.23).
- UM 461 - To get a better match of the IRS spectrum, we add an extra MIR modified blackbody ( $\beta$ fixed to 2.0 and best fit temperature $T=150 \mathrm{~K}$ ). This is shown in Fig. 7.19.


Fig. 7.19. SED for UM461 without (left) and with (right) the extra MIR modified blackbody.

### 7.2.5 Notes on individual KINGFISH galaxies

As we are interested in getting SEDs with Herschel constraints to have an accurate estimation of the dust mass with submm constraints, and also in the comparison of Spitzer-only versus Herschel dust masses, we do not consider some KINGFISH galaxies which have no Herschel detections: NGC 0584 - NGC 1404 - DDO 154 - DDO 165.

- NGC 0855 - If we leave the $f_{v s g}$ parameter free, this leads to unrealistically prominent PAH features for this galaxy. Thus we fix the $f_{v s g}$ parameter to the Galactic value for this galaxy. This leads to an increase of the dust mass of $\sim 38 \%$ which is within the $\sim 40 \%$ error bar. This is shown in Fig. 7.20.


Fig. 7.20. SEDs for NGC0855 with $\mathrm{f}_{\text {vsg }}$ free (left) and fixed (right).

- Ho $I$ - As for NGC 0855, we fix the $f_{v s g}$ parameter to the Galactic value for this galaxy. This leads to an increase of the dust mass of $\sim 14 \%$ which is within the $\sim 47 \%$ error bar.


Fig. 7.21. SEDs for NGC4254 (top) and NGC4579 (bottom) with $\mathrm{f}_{\text {vsg }}$ free (left) and fixed (right).

- NGC 1512 - As for NGC 0855 , we fix the $f_{v s g}$ parameter to the Galactic value for this galaxy. This leads to an increase of the dust mass of $\sim 17 \%$ which is within the $\sim 28 \%$ error bar.
- NGC 4254 and NGC 4579 - We fix the $f_{v s g}$ parameter to the Galactic value for these galaxies as it enables a better fit in the FIR. This is shown in Fig. 7.21. The dust mass for these two galaxies thus decreases by $\sim 20 \%$.
- NGC 5408 - As for NGC 0855 , we fix the $f_{v s g}$ parameter to the Galactic value for this galaxy. This leads to an increase of the dust mass of $\sim 16 \%$ which is within the $\sim 26 \%$ error bar.


### 7.2.6 Results



Fig. 7.22. Distributions of $M_{\text {dust }}$ (top left), $L_{T I R}$ (top right), $\langle U\rangle$ (bottom left) and $f_{P A H}$ (bottom right) from our SED fits for the DGS and KINGFISH samples. The colour scale represents the range of metallicity values. On each panel, the upper/lower histogram is the KINGFISH/DGS distribution for the parameter.

The individual SEDs for the DGS and KINGFISH galaxies are presented in Figs. 7.23 and 7.24. Note that in Figs. 7.23 and 7.24 the last fitted point is the $350 \mu \mathrm{~m}$ point, as explained in the beginning of Section 7.2. All longer wavelength points are just overlaid on the plots. The resulting parameters and their errors are given in Tables 7.3, 7.4, 7.5 and 7.6. The dust mass, TIR
luminosity, averaged starlight intensity and PAH mass fraction are presented in Fig. 7.22 and the main informations on these four parameters are summarised in Table 7.2.

The two samples span a range of 6 orders of magnitude in dust mass and in TIR luminosity. The mean starlight intensity ranges from 0.335 (Ho I) and $8.9 \times 10^{3}$ (Tol 1214-277). To determine the PAH mass fraction range, we consider only galaxies for which the $f_{v s g}$ parameter is free. For these galaxies, the PAH mass fraction ranges from 0.06 (for NGC 2366) to 1.6 (for IC 0342) times the Galactic value.

With these histograms we can already note a trend between all of the parameters and metallicity. The dust masses, TIR luminosities and PAH mass fractions are higher in the sample containing more metal-rich galaxies: the median $M_{\text {dust }}$ is $2.9 \times 10^{5} M_{\odot}$ for the DGS galaxies versus $1.6 \times 10^{7} M_{\odot}$ for the KINGFISH sample, the median $L_{T I R}$ is $9.3 \times 10^{8} L_{\odot}$ for the DGS galaxies versus $6.1 \times 10^{9} L_{\odot}$ for the KINGFISH sample, and $f_{P A H, D G S}^{m e d}=0.18 f_{P A H, \odot}$ compared to $f_{P A H, K I N G F I S H}^{m e d}=0.89$ $f_{P A H, \odot}$. This is consistent with the main dwarf galaxies characteristics we reviewed in Chapter 2: dwarf galaxies are less massive, less luminous and show weak PAH features. The mean starlight intensity behaves the opposite way as it decreases with increasing metallicity. This is consistent with dwarf galaxies harbouring a more intense galaxy wide radiation field.

Table 7.2. Main characteristics of the DGS and KINGFISH SED parameters.

|  | $\log \left(M_{\text {dust }}\right)$ <br> $\left[\log \left(M_{\odot}\right)\right]$ | $\log \left(L_{T I R}\right)$ <br> $\left[\log \left(L_{\odot}\right)\right]$ | $\log (<U>)$ | $f_{P A H}$ |
| :--- | :---: | :---: | :---: | :---: |
| DGS | $2.57-7.42$ | $6.50-11.30$ | $0.002-3.95$ | $0.00^{*}-0.54$ |
|  | median: 5.46 | median: 8.97 | median: 1.25 | median: 0.18 |
| KINGFISH | $4.02-8.02$ | $6.67-11.13$ | $-0.48-1.80$ | $0.00^{*}-1.59$ |
|  | median: 7.21 | median: 9.78 | median: 0.34 | median: 0.89 |

[^42]
### 7.2.7 SEDs fits for the DGS and KINGFISH samples


$\left[{ }^{\circ} \mathrm{T}\right]{ }^{n} \mathrm{~T} \wedge$

$\left[{ }^{\circ} \mathrm{T}\right]{ }^{\wedge} \mathrm{T} \cap$

1000
Fig. 7.23. SEDs for Haro11 (top left), Haro2 (top right), Haro3 (bottom left), He2-10 (bottom right). The observed SED includes the Herschel data (purple crosses) as well as any and
for IRAS, orange crosses for SCUBA or LABOCA, squares for MAMBO. The IRS spectrum is displayed in dark blue, with the fitted parts in orange. The total modelled SED in black is the sum of the stellar (green) and dust (red) contributions, and eventually an extra MIR modified blackbody contribution (yellow). The modelled points in the different bands are the filled blue circles.



Fig. 7.23. (continued) for HS0017+1055 (top left), HS0052+2536 (top right), HS0822+3542 (bottom left), HS1222+3741 (bottom right).

$10 \underset{\text { Wavelength }}{[\mu \mathrm{m}]} \mathrm{CO}$
Wavelength $[\mu \mathrm{m}]$





$\left[{ }^{\circ} \mathrm{T}\right]{ }^{n} \mathrm{~T} /$



$$
1000
$$

Wavelength [ $\mu \mathrm{m}$ ]
Fig. 7.23. (continued) for Mrk153 (top left), Mrk209 (top right), Mrk930 (bottom left), NGC1140 (bottom right).

$\left[{ }^{\circ} \mathrm{T}\right]{ }^{n} \mathrm{~T} \wedge$

1000

$\left[{ }^{\circ} \mathrm{T}\right]{ }^{n} \mathrm{~T} \Lambda$


Wavelength [ $\mu \mathrm{m}$ ]
Fig. 7.23. (continued) for NGC1569 (top left), NGC1705 (top right), NGC2366 (bottom left), NGC4214 (bottom right).






Fig. 7.23. (continued) for SBS1211+540 (top left), SBS1249+493 (top right), SBS1415+437 (bottom left), SBS1533+574 (bottom right).


## d) for SBS1211 +50



1000
$\left[{ }^{\circ} \mathrm{T}\right]{ }^{n} \mathrm{~T} \Omega$
sןenpisəəy


$\left[{ }^{\circ} \mathrm{T}\right]{ }^{n} \mathrm{~T} \wedge$


[^43]




Fig. 7.23. (continued) for UM448 (top left), UM461 (top right), VIIZw403 (bottom left).

Fig. 7.24. Same as Fig. 7.23 for NGC0337 (top left), NGC0628 (top right), NGC0855 (bottom left), NGC0925 (bottom right).




Fig. 7.24. (continued) for NGC1097 (top left), NGC1266 (top right), NGC1291 (bottom left), NGC1316 (bottom right).
 Wavelength $[\mu \mathrm{m}]$

Fig. 7.24. (continued) for NGC1377 (top left), IC0342 (top right), NGC1482 (bottom left), NGC1512 (bottom right).

8

NGC2798

Fig. 7.24. (continued) for NGC2146 (top left), HoII (top right), DDO053 (bottom left), NGC2798 (bottom right).

000100 I 0 I DDO053


101001000

Fig. 7.24. (continued) for NGC2841 (top left), NGC2915 (top right), HoI (bottom left), NGC2976 (bottom right).




${ }^{10} \begin{array}{ll}\text { Wavelength }[\mu \mathrm{m}] & 1000\end{array}$
Fig. 7.24. (continued) for NGC3049 (top left), NGC3077 (top right), M81dwB (bottom left), NGC3190 (bottom right).




Fig. 7.24. (continued) for NGC3184 (top left), NGC3198 (top right), IC2574 (bottom left), NGC3265 (bottom right).



Fig. 7.24. (continued) for NGC3351 (top left), NGC3521 (top right), NGC3621 (bottom left), NGC3627 (bottom right).


Wavelength $[\mu \mathrm{m}$ ]
NGC4254

Fig. 7.24. (continued) for NGC3773 (top left), NGC3938 (top right), NGC4236 (bottom left), NGC4254 (bottom right).


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Fig. 7.24. (continued) for NGC4321 (top left), NGC4536 (top right), NGC4559 (bottom left), NGC4569 (bottom right).

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Fig. 7.24. (continued) for NGC4579 (top left), NGC4594 (top right), NGC4625 (bottom left), NGC4631 (bottom right).

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Wavelength [ $\mu \mathrm{m}$ ]
Fig. 7.24. (continued) for NGC4725 (top left), NGC4736 (top right), NGC4826 (bottom left), NGC5055 (bottom right).



Fig. 7.24. (continued) for NGC5398 (top left), NGC5408 (top right), NGC5457 (bottom left), NGC5474 (bottom right).


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 10 Wavelength $[\mu \mathrm{m}] \quad 1000$
Fig. 7.24. (continued) for NGC5713 (top left), NGC5866 (top right), NGC6946 (bottom left), NGC7331 (bottom right).

Fig. 7.24. (continued) for $\mathrm{NGC7793}$ (top left).
Table 7.3. SED parameters: $M_{d u s t}, L_{T I R}, \mathrm{SFR},\langle U\rangle, U_{\min }, \Delta U$ and $\alpha$ for the DGS sample.

| Source | $\begin{aligned} & \hline M_{\text {dust }} \\ & {\left[M_{\odot}\right]} \end{aligned}$ | $\begin{aligned} & \hline L_{T I R} \\ & {\left[L_{\odot}\right]} \end{aligned}$ | $\begin{gathered} \text { SFR } \\ {\left[\mathrm{M}_{\odot} \mathrm{yr}^{-1}\right]} \end{gathered}$ | $<U>$ | $U_{\text {min }}$ | $\Delta U$ | $\alpha$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Haro11 | $9.9{ }_{-0 .}^{+1.0} \times 10^{6}$ | $2.0_{-0.04}^{+0.04} \times 10^{11}$ | $3.4{ }_{-0.07}^{+0.08} \times 10^{1}$ | $1.4_{-0.2}^{+0.2} \times 10^{2}$ | $11.60{ }_{-2.6}^{+2.24}$ | $3.6_{-16}^{+3.4} \times 10^{5}$ | $1.96{ }_{-0.02}^{+0.01}$ |
| Haro2 | $1.7_{-0.2}^{+0.2} \times 10^{6}$ | $6.4_{-0.14}^{+0.13} \times 10^{9}$ | $1.1_{-0.02}^{+0.02}$ | $2.3_{-0.2}^{+0.3} \times 10^{1}$ | $3.03_{-0.4}^{+0.50}$ | $7.3_{-2.0}^{+2.0} \times 10^{5}$ | $2.10_{-0.01}^{+0.01}$ |
| Haro3 | $2.3_{-0.2}^{+0.3} \times 10^{6}$ | $5.4_{-0.16}^{+0.13} \times 10^{9}$ | $9.33_{-0.28}^{+0.22} \times 10^{-1}$ | $1.55_{-0.2}^{+0.2} \times 10^{1}$ | $2.03_{-0.4}^{+0.36}$ | $6.3_{-1.8}^{+6.3} \times 10^{5}$ | $2.11_{-0.02}^{+0.02}$ |
| He2-10 | $1.6_{-0.2}^{+0.2} \times 10^{6}$ | $5.5{ }_{-0.12}^{+0.13} \times 10^{9}$ | $9.55_{-0.22}^{+0.22} \times 10^{-1}$ | $2.3_{-0.2}^{+0.3} \times 10^{1}$ | $2.211_{-0.3}^{+0.45}$ | $3.7_{-0.5}^{+0.6} \times 10^{5}$ | $2.03_{-0.01}^{+0.02}$ |
| HS0017+1055 | $8.3_{-2.8}^{+7.9} \times 10^{4}$ | $9.7_{-0.49}^{-0.55} \times 10^{8}$ | $1.7_{-0.08}^{+0.09} \times 10^{-1}$ | $8.0_{-4.5}^{+5.0} \times 10^{1}$ | $5.14_{-3.6}^{+4.52}$ | $2.9_{-0.5}^{+0.5} \times 10^{6}$ | $1.97_{-0.04}^{+0.03}$ |
| HS0052+2536 | $6.6_{-2.1}^{+2.4} \times 10^{6}$ | $1.7_{-0.08}^{+0.08} \times 10^{10}$ | ${ }^{2}$ | $1.7_{-0.4}^{+0.8} \times 10^{1}$ | $2.90{ }_{-1}^{+1 .}$ | $3.4_{-1.9}^{+3.0} \times 10^{5}$ | $2.18{ }_{-0.05}^{+0.04}$ |
| HS0822+3542 | $6.2_{-1.9}^{+4.2} \times 10^{3}$ | $1.11_{-0.09}^{+0.10} \times 10^{7}$ | $1.8{ }_{-0.15}^{+0.18} \times 10^{-3}$ | $1.11_{-0.5}^{+0.5} \times 10^{1}$ | $1.92{ }_{-1}^{+1 .}$ | $1.1_{-0.2}^{+0.5} \times 10^{5}$ | $2.17_{-0.06}^{+0.06}$ |
| HS1222+3741 | $8.0_{-2.6}^{+2.9} \times 10^{5}$ | $2.4{ }_{-0.24}^{+0.17} \times 10^{9}$ | $4.22_{-0.42}^{+0.30} \times 10^{-1}$ | $1.9{ }_{-0.5}^{+0.9} \times 10^{1}$ | $1.96{ }_{-0.5}^{+1.07}$ | $9.3{ }_{-2.4}^{+3.6} \times 10^{5}$ | $2.05_{-0.03}^{+0.03}$ |
| HS1236+3937 |  |  |  |  |  |  |  |
| HS1304+3529 | $4.0_{-2.2}^{+2.5} \times 10^{6}$ | $1.7_{-0.07}^{+0.09} \times 10^{9}$ | $3.0_{-0.13}^{+0.15} \times 10^{-1}$ | $2.6_{-1.6}^{+15.4}$ | $0.01_{-0.0}^{+3.90}$ | $3.0_{-0 .}^{+12.5} \times 10^{2}$ | $1.44_{-0.16}^{+0.67}$ |
| HS1319+3224 | $2.9{ }_{-2.3}^{+14.8} \times 10^{4}$ | $2.4{ }_{-0.26}^{+0.24} \times 10^{8}$ | $4.1_{-0.45}^{+0.42} \times 10^{-2}$ | $4.2{ }_{-3.4}^{+5.9} \times 10^{1}$ | $7.73_{-4.8}^{+0.10}$ | $1.22_{-8.8}^{+33.8} \times 10^{5}$ | $2.18{ }_{-0.09}^{+0.08}$ |
| HS1330+3651 | $1.3_{-0.2}^{+0.3} \times 10^{6}$ | $1.4{ }_{-0.06}^{+0.06} \times 10^{9}$ | $2.4{ }_{-0.10}^{+0.10} \times 10^{-1}$ | $6.9_{-1.5}^{+1.7}$ | $1.31_{-0.3}^{+0.39}$ | $2.5{ }_{-1.3}^{+3.9} \times 10^{4}$ | $2.19_{-0.04}^{+0.03}$ |
| HS1442+4250 | - | - |  |  |  |  |  |
| HS2352+2733 | - | - | - | - | - |  | - |
| IZw18 | $5.1_{-0.7}^{+1.1} \times 10^{2}$ | $3.3{ }_{-0.19}^{+0.20} \times 10^{7}$ | $5.8{ }_{-0.33}^{+0.34} \times 10^{-}$ | $3.9{ }_{-0.8}^{+0.3} \times 10^{2}$ | $383.54_{-107.9}^{+24.81}$ | $2.1_{-1.6}^{+7.6} \times 10^{1}$ | $1.00{ }_{-0.00}^{+0.001}$ |
| IC10 | $2.8{ }_{-0.8}^{+1.1} \times 10^{5}$ | $1.7_{-0.07}^{+0.06} \times 10^{8}$ | $2.9{ }_{-0.13}^{+0.11} \times 10^{-2}$ | $3.9{ }_{-1.1}^{+1.4}$ | $0.72_{-0.2}^{+0.30}$ | $1.0_{-0.2}^{+0.2} \times 10^{6}$ | $2.21_{-0.02}^{+0.02}$ |
| IIZw40 | $6.7_{-1.6}^{+4.2} \times 10^{5}$ | $2.9{ }_{-0.16}^{+0.10} \times 10^{9}$ | $5.0_{-0.27}^{+0.17} \times 10^{-1}$ | $2.3{ }_{-1.0}^{+0.8} \times 10^{1}$ | $1.988_{-1.4}^{+2.00}$ | $1.3{ }_{-0.9}^{+4.3} \times 10^{4}$ | $1.92_{-0.29}^{+0.16}$ |
| Mrk1089 | $2.6_{-0.3}^{+0.4} \times 10^{7}$ | $3.6{ }_{-0.10}^{+0.10} \times 10^{10}$ | ${ }_{6.2} 2_{-0.17}^{+0.17}$ | $8.6_{-1.1}^{+1.4}$ | $1.45{ }_{-0.3}^{\text {-0. }}$. ${ }^{\text {a }}$ | $6.8{ }_{-1.9}^{+2.9} \times 10^{4}$ | $2.16{ }_{-0.03}^{+0.03}$ |
| Mrk1450 | $5.0_{-1.9}^{+18.9} \times 10^{4}$ | $3.0{ }_{-0.08}^{+0.10} \times 10^{8}$ | $5.2_{-0.14}^{+0.16} \times 10^{-2}$ | $3.1_{-2.7}^{+2.0} \times 10^{1}$ | $1.88_{-2.6}^{+6.78}$ | $1.1_{-0.4}^{+7.0} \times 10^{3}$ | $1.63_{-0.21}^{+0.44}$ |
| Mrk153 | $1.7_{-0.3}^{+0.4} \times 10^{5}$ | $1.0_{-0.04}^{+0.04} \times 10^{9}$ | $1.88_{-0.06}^{+0.07} \times 10^{-1}$ | $7.0_{-1.6}^{+1.9} \times 10^{1}$ | $15.48_{-4.8}^{+6.49}$ | $1.1_{-0.3}^{+0.6} \times 10^{5}$ | $2.25-0.05$ |
| Mrk209 | $4.6_{-1.4}^{+2.1} \times 10^{3}$ | $2.8{ }_{-0.13}^{+0.16} \times 10^{7}$ | $4.9{ }_{-0.23}^{+0.27} \times 10^{-3}$ | $4.1_{-1.4}^{+2.0} \times 10^{1}$ | $5.28_{-2.5}^{+4.09}$ | $1.9_{-0.8}^{+1.2} \times 10^{5}$ | $2.07_{-0.06}^{+0.06}$ |
| Mrk930 | $6.0_{-5.6}^{+0.9} \times 10^{6}$ | $1.9{ }_{-0.08}^{+0.13} \times 10^{10}$ | $3.3{ }_{-0.14}^{+0.22}$ | $1.9{ }_{-0.7}^{+0.3} \times 10^{1}$ | $3.18_{-0.6}^{+0.56}$ | $3.6{ }_{-1.9}^{+0.3} \times 10^{4}$ | $2.15{ }_{-0.04}^{+0.34}$ |
| NGC1140 | $3.9{ }_{-0.5}^{+0.5} \times 10^{6}$ | $3.7_{-0.12}^{+0.13} \times 10^{9}$ | $6.5{ }_{-0.22}^{+0.23} \times 10^{-1}$ | $6.1_{-0.8}^{+0.8}$ | $1.07_{-0.2}^{-0.62}$ | $7.7_{-2.1}^{+4.5} \times 10^{4}$ | $2.188_{-0.02}^{+0.02}$ |
| NGC1569 | $2.9{ }_{-0.4}^{+0.5} \times 10^{5}$ | $1.3{ }_{-0.03}^{+0.03} \times 10^{9}$ | $2.22_{-0.06}^{+0.05} \times 10^{-1}$ | $2.4{ }_{-0.4}^{+0.4} \times 10^{1}$ | $5.53_{-1.9}^{+1.46}$ | $1.1_{-0.7}^{+0.8} \times 10^{4}$ | $2.23_{-0.13}^{+0.06}$ |
| NGC1705 | $1.3{ }_{-0.7}^{+0.8} \times 10^{5}$ | $6.9_{-0.39}^{+0.31} \times 10^{7}$ | $1.22_{-0.07}^{+0.05} \times 10^{-2}$ | $3.6{ }_{-1.6}^{+4.1}$ | $0.18_{-0.2}^{+1.58}$ | $4.66_{-13}^{+230.1} \times 10^{2}$ | $1.73_{-0.10}^{+0.52}$ |
| NGC2366 | $1.2_{-0.3}^{+0.5} \times 10^{5}$ | $1.4{ }_{-0.06}^{+0.06} \times 10^{8}$ | $2.44_{-0.10}^{+0.10} \times 10^{-2}$ | $8.0_{-2.3}^{+2.8}$ | $0.61{ }_{-0.3}^{+0.62}$ | $1.2_{-0.5}^{+1.9} \times 10^{4}$ | $1.933_{-08}^{+0.11}$ |
| NGC4214 | $1.7_{-0.4}^{+0.5} \times 10^{6}$ | $5.4_{-0.15}^{+0.16} \times 10^{8}$ | $9.44_{-0.27}^{+0.28} \times 10^{-2}$ | $2.0_{-0.5}^{+0.6}$ | $0.31_{-0.1}^{+0.15}$ | $3.3_{-3.3}^{+7.4} \times 10^{3}$ | $2.10_{-0.06}^{+0.07}$ |
| NGC4449 | $3.0_{-0.3}^{+0.3} \times 10^{6}$ | $2.22_{-0.07}^{+0.06} \times 10^{9}$ | $3.8{ }_{-0.11}^{+0.10} \times 10^{-1}$ | $4.6{ }_{-0.5}^{+0.6}$ | $1.05_{-0.1}^{-0.17}$ | $4.7_{-1.1}^{+1.1} \times 10^{4}$ | $2.288_{-0.02}^{+0.02}$ |
| NGC4861 | $6.0_{-1.2}^{+1.9} \times 10^{5}$ | $3.5{ }_{-0.12}^{+0.11} \times 10^{8}$ | $6.0_{-0.21}^{+0.20} \times 10^{-2}$ | $3.8{ }_{-0.9}^{+1.1}$ | $0.20_{-0.1}^{+0.09}$ | $1.6{ }_{-0.5}^{+0.9} \times 10^{4}$ | $1.900_{-0.04}^{+0.03}$ |
| NGC5253 | $5.6_{-0.9}^{+1.6} \times 10^{5}$ | $1.6_{-0.03}^{+0.03} \times 10^{9}$ | $2.8{ }_{-0.05}^{+0.05} \times 10^{-1}$ | $1.4{ }_{-0.3}^{+0.3} \times 10^{1}$ | $2.38{ }_{-1.2}^{-1.03}$ | $5.6{ }_{-3.8}^{+10.4} \times 10^{3}$ | $2.12_{-0.18}^{+0.10}$ |
| NGC625 | $4.5{ }_{-0.7}^{+0.8} \times 10^{5}$ | $2.9{ }_{-0.08}^{+0.07} \times 10^{8}$ | $5.0_{-0.15}^{+0.12} \times 10^{-2}$ | $4.1_{-0.7}^{+0.8}$ | $0.55_{-0.1}^{+0.13}$ | $7.0_{-4.2}^{+6.8} \times 10^{5}$ | $2.12_{-0.03}^{+0.03}$ |
| NGC6822 | $2.3_{-1.5}^{+3.8} \times 10^{5}$ | $3.66_{-0.18}^{+0.20} \times 10^{7}$ | $6.3_{-0.31}^{+0.35} \times 10^{-3}$ | $1.0_{-0.6}^{+1.1}$ | $0.08_{-0.1}^{+0.18}$ | $3.9_{-1.9}^{+4.3} \times 10^{3}$ | $1.97_{-0.09}^{+0.09}$ |

Table 7.3. SED parameters: $M_{d u s t}, L_{T I R}, \mathrm{SFR},\langle U\rangle, U_{\min }, \Delta U$ and $\alpha$ for the DGS sample.

| rce | $\begin{aligned} & M_{\text {dust }} \\ & {\left[M_{\odot}\right]} \end{aligned}$ | $\begin{gathered} \hline L_{T I R} \\ {\left[L_{\odot}\right]} \end{gathered}$ | $\begin{gathered} \text { SFR } \\ {\left[\mathrm{M}_{\odot} \mathrm{yr}^{-1}\right]} \end{gathered}$ | <U> | $U_{\text {min }}$ | $\Delta U$ | ${ }^{\alpha}$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Pox186 | $1.9{ }_{-0.7}^{+1.9} \times 10^{5}$ | $5.7_{-0.37}^{+0.31} \times 10^{7}$ | $9.9 .9 .0 .64 \times 10^{-3}$ | $3.44_{-1.2}^{+8.7} \times 10^{1}$ | $3.44+$ | $4.6{ }_{-0.1}^{+0.4} \times 10^{4}$ | $1.99_{-0.71}^{+0.28}$ |
| 35-052 | $3.1_{-4.5}^{+160.0} \times 10^{3}$ | $1.8 .8{ }_{-0.06}^{+0.03} \times 10^{9}$ | $3.0{ }_{-0.11}^{+0.06} \times 10^{-1}$ | $1.11_{0.3}^{+0.3} \times 10^{3}$ | 1070.37 | $1.0{ }_{-3248.6}^{+1429.7}$ | $1.00_{-0.00}^{+0.69}$ |
| SBS1159+545 | $3.0{ }_{-2.1}^{+6.8} \times 10^{4}$ | $2.0_{-0.10}^{+0.09} \times 10^{8}$ | $3.5{ }_{-0.16}^{+0.15} \times 10^{-2}$ | $4.66_{-2.4}^{+3.3} \times 10^{1}$ | $3.33+$ | $1.2{ }_{-0.3}^{+0.3} \times 10^{6}$ | $1.99_{-0.04}^{+0.05}$ |
| SBS1211+540 | $1.88_{-0.9}^{+0.9} \times 10^{4}$ | $2.9{ }_{-0.15}^{+0.17} \times 10^{7}$ | $4.9 .9 .0 .26 \times 10^{-3}$ | $1.00_{-0.4}^{+1.2} \times 10^{1}$ | $0.02+$ | $1.4{ }_{-0.3}^{+0.3} \times 10^{3}$ | $1.42_{-0.11}^{+0.088}$ |
| SBS1249+493 | $4.22_{-1.0}^{+1.7} \times 10^{5}$ | $9.33_{-0.71}^{+0.88} \times 10^{8}$ | $1.6_{-0.12}^{+0.15} \times 10^{-}$ | $1.44_{-0.7}^{+0.6} \times 10^{1}$ | 2.30 | $4.8{ }_{-0.8}^{\text {Pr. }} \times 10^{5}$ | $2.16_{-0.03}^{+0.04}$ |
| SBS1415+437 | $6.44_{-1.1}^{1.2} \times 10^{3}$ | $6.5{ }_{-0.31}^{+0.27} \times 10^{7}$ | $1.11_{-0.05}^{+0.05} \times 10^{-2}$ | $6.99_{-1.0}^{+1.4} \times 10^{1}$ | 63 | $3.2{ }_{-0.7}^{+1.0} \times 10^{5}$ | 2.24 |
| SBS1533+574 | $2.0{ }_{-0.4}^{+0.5} \times 10^{6}$ | $1.9{ }_{-0.07}^{+0.05} \times 10^{9}$ | $3.33_{-0.13}^{+0.09} \times 10^{-1}$ | 4.4 | 0.67 | $6.0{ }_{-2.5}^{+2.9} \times 10^{3}$ | 2.10 |
| Tol0618-402 |  |  |  |  |  |  |  |
| Tol1214-277 | $1.33_{-0.7}^{+2.3} \times 10^{5}$ | $8.88_{-0.49}^{+0.40} \times 10^{8}$ | $1.5{ }_{-0.08}^{+0.07} \times 10^{-1}$ | $8.9_{-0.6}^{+4.5} \times 10^{3}$ | $0.89{ }_{-0}^{+1}$ | $1.0_{-0.2}^{+0.9} \times 10^{5}$ | 1.00 |
| UGC4483 | $3.3{ }_{-1.4}^{+1.7} \times 10^{2}$ | $3.1{ }_{-0.25}^{+0.25} \times 10^{6}$ | $5.44_{-0.44}^{+0.44} \times 10^{-4}$ | $5.44_{-1.7}^{+2.9} \times 10^{1}$ | $18.27_{-5.9}^{+8.03}$ | $4.6{ }_{-3.3}^{+14.6} \times 10^{4}$ | $2.50{ }_{-1.46}^{+0.04}$ |
| UGCA20 |  |  |  |  |  |  |  |
| UM133 | $9.5-3.9 \times 10^{4}$ | $7.2_{-0.59}^{+0.65} \times 10^{7}$ | $1.22_{-0.10}^{+0.11} \times 10$ | $5.1+$ | $0.50{ }_{-0}^{+0}$ | $1.8{ }_{-0.7}^{+0.6} \times 10^{5}$ | $2.04{ }_{-0.04}^{+0.05}$ |
| UM311 | $2.00_{-0.3}^{+0.4} \times 10^{7}$ | $5.5_{-0.20}^{+0.21} \times 10^{9}$ | $9.4{ }_{-035}^{+0.37} \times 10^{-1}$ | $1.6{ }_{-0.3}^{+0.3}$ | $0.38_{-0.1}^{+0.11}$ | $1.0{ }_{-0.4}^{+0.0} \times 10^{7}$ | 2.30 |
| UM448 | $1.8{ }_{-0.2}^{0+1} \times 10^{7}$ | $8.9_{-0.26}^{+0.22} \times 10^{10}$ | $1.5{ }_{-0.04}^{+0.04} \times 10^{1}$ | $3.22_{-0.2}^{+0.3} \times 10^{1}$ | $5.66{ }_{-0.7}^{+0.72}$ | $5.7{ }_{-2.1}^{-0.6 .6} \times 10^{5}$ | $2.19+$ |
| UM | $2.11_{-0.7}^{+1.4} \times 10^{4}$ | $7.1{ }_{-0.21}^{+0.25} \times 10^{7}$ | $1.2{ }_{-0.04}^{+0.04} \times 10^{-2}$ | $2.00_{0.9}^{+0.9} \times 10^{1}$ | $1.79_{-1}^{+1}$ | $1.3{ }_{-0.3}^{+0.3} \times 10^{5}$ | $2.00_{-0.04}^{+0.03}$ |
| VIIZw403 | $7.22_{-1.1}^{+1.3} \times 10^{3}$ | $2.0{ }_{-0.07}^{0.07} \times 10^{7}$ | $3.5_{-0.12}^{+0.12} \times 10^{-3}$ | $1.88_{-0.3}^{+0.9} \times 10^{1}$ | $4.29_{-1.0}^{+1.15}$ | $3.6{ }_{-1.9}^{+3.9} \times 10^{4}$ | $2.28_{-0.05}^{+0.04}$ |

Table 7.4. SED parameters: $f_{v s g}, f_{P A H}, f_{i o n}$ and $M_{\star}$ for the DGS sample.

| Source | $f_{v s g}$ | $f_{P A H}$ | $f_{\text {ion }}$ | $\begin{gathered} M_{\star} \\ {\left[M_{\odot}\right]} \\ \hline \end{gathered}$ |
| :---: | :---: | :---: | :---: | :---: |
| Haro11 | $1.63_{-0.38}^{+0.22}$ | $0.09_{-0.01}^{+0.01}$ | 1.00 | - |
| Haro2 | $0.25_{-0.19}^{+0.19}$ | $0.54_{-0.02}^{+0.02}$ | 1.00 | $7.7_{-0.23}^{+0.23} \times 10^{9}$ |
| Haro3 | $0.23_{-0.12}^{+0.28}$ | $0.36_{-0.02}^{+0.02}$ | $0.22_{-0.05}^{+0.04}$ | $6.0_{-0.16}^{+0.19} \times 10^{9}$ |
| He2-10 | $0.39_{-0.09}^{+0.09}$ | $0.48_{-0.02}^{+0.03}$ | $0.46_{-0.04}^{+0.05}$ | $4.9_{-0.12}^{+0.12} \times 10^{9}$ |
| HS0017+1055 | 0.00 | 0.00 | 0.50 | $8.6_{-0.34}^{+0.34} \times 10^{8}$ |
| HS0052+2536 | $0.37_{-0.22}^{+0.21}$ | $0.18_{-0.07}^{+0.07}$ | 0.50 | $6.2_{-0.23}^{+0.25} \times 10^{9}$ |
| HS0822+3542 | $0.31_{-0.09}^{+0.10}$ | 0.00 | 0.50 | $8.3_{-0.66}^{+0.60} \times 10^{6}$ |
| HS1222+3741 | 1.00 | 0.00 | 0.50 | $7.4_{-0.37}^{+0.33} \times 10^{9}$ |
| HS1236+3937 | - | - | - |  |
| HS1304+3529 | $1.06_{-0.24}^{+0.20}$ | 0.00 | 0.50 | $6.1_{-0.27}^{+0.27} \times 10^{8}$ |
| HS1319+3224 | $0.90_{-0.40}^{+0.91}$ | 0.00 | 0.50 | $2.5_{-1.40}^{+0.82} \times 10^{8}$ |
| HS1330+3651 | $1.22_{-0.19}^{+0.17}$ | 0.00 | 0.50 | $2.0_{-0.48}^{+0.45} \times 10^{9}$ |
| HS1442+4250 | - | - | - | - |
| HS2352+2733 | - | - | - | - |
| IZw18 | $2.07_{-0.30}^{+0.29}$ | 0.00 | 0.50 | $7.0_{-0.45}^{+0.58} \times 10^{7}$ |
| IC10 | $0.00_{-0.01}^{+0.01}$ | $0.29_{-0.01}^{+0.01}$ | $0.22_{-0.01}^{+0.01}$ | $4.3_{-0.06}^{+0.07} \times 10^{8}$ |
| IIZw40 | $0.99_{-0.33}^{+0.39}$ | $0.17_{-0.04}^{+0.05}$ | 0.50 | $8.6_{-0.59}^{+0.61} \times 10^{8}$ |
| Mrk1089 | $1.23_{-0.24}^{+0.22}$ | $0.18_{-0.07}^{+0.07}$ | $0.70_{-0.42}^{+0.28}$ | $1.3{ }^{+0.04}+10.040^{10}$ |
| Mrk1450 | $0.66_{-0.26}^{+0.19}$ | $0.12_{-0.02}^{+0.01}$ | $0.04_{-0.05}^{+0.19}$ | $1.22_{-0.03}^{+0.03} \times 10^{8}$ |
| Mrk153 | $0.933_{-0.09}^{+0.08}$ | 0.00 | 0.50 | $2.6_{-0.17}^{+0.20} \times 10^{9}$ |
| Mrk209 | $0.26_{-0.24}^{+0.28}$ | 0.00 | 0.50 | $8.7_{-1.06}^{+0.88} \times 10^{7}$ |
| Mrk930 | $0.52_{-0.12}^{+0.28}$ | $0.07_{-0.02}^{+0.03}$ | 1.00 | $8.1_{-0.55}^{+0.79} \times 10^{9}$ |
| NGC1140 | $0.66_{-0.12}^{+0.12}$ | $0.47_{-0.02}^{+0.02}$ | $0.44_{-0.03}^{+0.03}$ | $7.1_{-0.21}^{+0.23} \times 10^{9}$ |
| NGC1569 | $0.70_{-0.03}^{+0.03}$ | $0.17_{-0.01}^{+0.01}$ | $0.25_{-0.03}^{+0.03}$ | $1.8_{-0.10}^{+0.28} \times 10^{9}$ |
| NGC1705 | $0.60_{-0.58}^{+0.08}$ | $0.07_{-0.05}^{+0.12}$ | 1.00 | $4.7_{-0.28}^{+0.27} \times 10^{8}$ |
| NGC2366 | $0.46_{-0.33}^{+0.57}$ | $0.06_{-0.08}^{+0.07}$ | $1.00_{-0.94}^{+0.00}$ | $4.8_{-0.32}^{+0.29} \times 10^{8}$ |
| NGC4214 | $1.52_{-0.52}^{+0.45}$ | $0.27_{-0.12}^{+0.16}$ | $0.89_{-0.70}^{+0.05}$ | $1.6{ }_{-0.05}^{+0.05} \times 10^{9}$ |
| NGC4449 | $0.57_{-0.03}^{+0.04}$ | $0.54_{-0.02}^{+0.02}$ | $0.78_{-0.02}^{+0.01}$ | $5.9_{-0.26}^{+0.22} \times 10^{9}$ |
| NGC4861 | $0.911_{-0.43}^{+0.20}$ | $0.00_{-0.01}^{+0.10}$ | 0.00 | $6.8{ }_{-1.20}^{+1.17} \times 10^{8}$ |
| NGC5253 | $2.40_{-0.12}^{+0.11}$ | $0.24_{-0.05}^{+0.07}$ | $0.57_{-0.30}^{+0.49}$ | $2.0_{-0.06}^{+0.07} \times 10^{9}$ |
| NGC625 | $0.76_{-0.60}^{+0.63}$ | $0.14_{-0.10}^{+0.13}$ | $0.711_{-0.50}^{+0.50}$ | $1.2{ }_{-0.04}^{+0.04} \times 10^{9}$ |
| NGC6822 | 1.00 | $0.12_{-0.06}^{+0.08}$ | $1.00_{-1.00}^{+0.00}$ | $3.4{ }_{-0.40}^{+0.32} \times 10^{8}$ |
| Pox186 | $0.46_{-0.72}^{+0.45}$ | 0.00 | 0.50 | $3.9_{-1.15}^{+0.92} \times 10^{7}$ |
| SBS0335-052 | $3.07_{-0.78}^{+0.88}$ | 0.00 | 0.50 | $4.5_{-0.25}^{+0.19} \times 10^{8}$ |
| SBS1159+545 | $0.00_{-0.05}^{+0.64}$ | 0.00 | 0.50 | $2.7_{-0.80}^{+0.255} \times 10^{8}$ |
| SBS1211+540 | $0.69_{-0.06}^{+0.06}$ | 0.00 | 0.50 | $2.22_{-0.47}^{+0.46} \times 10^{7}$ |
| SBS1249+493 | 1.00 | 0.00 | 0.50 | $2.0_{-1.96}^{+1.96} \times 10^{8}$ |
| SBS1415+437 | 0.00 | 0.00 | 0.50 | $1.8_{-0.20}^{+0.26} \times 10^{8}$ |
| SBS1533+574 | $1.84_{-0.14}^{+0.14}$ | $0.17_{-0.01}^{+0.01}$ | $0.14_{-0.07}^{+0.06}$ | $2.0_{-0.06}^{+0.06} \times 10^{9}$ |
| Tol0618-402 | - | - | - | - |
| Tol1214-277 | 1.00 | 0.00 | 0.50 | $5.0 \times 10^{8}$ |
| UGC4483 | $0.79_{-0.22}^{+0.20}$ | 0.00 | 0.50 | $1.0_{-0.08}^{+0.08} \times 10^{7}$ |
| UGCA20 | - | - | - | $-$ |
| UM133 | $0.21_{-0.18}^{+0.32}$ | 0.00 | 0.50 | $1.6_{-0.33}^{+0.24} \times 10^{8}$ |
| UM311 | 1.00 | $0.78_{-0.19}^{+0.22}$ | 0.00 | $1.5_{-0.33}^{+0.31} \times 10^{10}$ |
| UM448 | $0.63_{-0.20}^{+0.20}$ | $0.27_{-0.09}^{+0.09}$ | $0.36_{-0.29}^{+0.34}$ | $5.4_{-0.16}^{+0.14} \times 10^{10}$ |
| UM461 | $0.55_{-0.16}^{+0.14}$ | 0.00 | 0.50 | $9.5{ }_{-0.55}^{+0.53} \times 10^{7}$ |
| VIIZw403 | $0.85_{-0.12}^{+0.13}$ | 0.00 | 0.50 | $2.7_{-0.14}^{+0.16} \times 10^{7}$ |

When no errors are given, it means that either this parameter was fixed in the fit or that the parameter reached the minimum or maximum value allowed for this parameter. If fixed: $f_{v s g}=1, f_{P A H}=0.0, f_{i o n}=0.5$, and $f_{v s g} \geq 0.0, f_{P A H} \geq 0.0$ and $f_{\text {ion }} \in[0.0,1.0]$.
Table 7.5. SED parameters: $M_{d u s t}, L_{T I R}, \mathrm{SFR},\langle U\rangle, U_{m i n}, \Delta U$ and $\alpha$ for the KINGFISH sample.

| Source | $\begin{aligned} & \hline M_{\text {dust }} \\ & {\left[M_{\odot}\right]} \end{aligned}$ | $\begin{aligned} & \hline L_{T I R} \\ & {\left[L_{\odot}\right]} \\ & \hline \end{aligned}$ | $\begin{gathered} \text { SFR } \\ {\left[\mathrm{M}_{\odot} \mathrm{yr}^{-1}\right]} \end{gathered}$ | $<U>$ | $U_{\text {min }}$ | $\Delta U$ | ${ }^{\alpha}$ |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| 7 | $1.6{ }_{-0.1}^{+0.2} \times 10^{7}$ | $1.1_{-0.03}^{+0.03} \times 10^{10}$ | $2.0{ }_{-0.05}^{+0.05}$ | $4.5{ }_{-0.4}^{+0.3} \times 10^{1}$ | $1.22_{-0.2}^{+0.18}$ | $8.9{ }_{-4.7}^{+1.8} \times 10^{5}$ | $2.3{ }_{-0.03}^{+0.03}$ |
| NGC0584 |  |  |  |  |  |  |  |
| NGC0628 | $2.66_{-0.3}^{+0.3} \times 10^{7}$ | $6.5{ }_{-0.17}^{+0.15} \times 10^{9}$ | $1.1{ }_{-0.03}^{+0.03}$ | $1.4_{-0.2}^{+0.2} \times 10^{1}$ | 0.57 | $1.8{ }_{-0.9}^{+2.2} \times 10^{1}$ | $50_{-0.00}^{+0.00}$ |
| NG | $1.85{ }_{-0.4}^{+0.5} \times 10^{6}$ | $3.9 .9{ }_{-0.10}^{+0.09} \times 10^{8}$ | $6.88_{-0.17}^{+0.15} \times 10^{-2}$ | $1.44_{-0.3}^{+0.3} \times 10^{1}$ | $0.09_{-0.03}^{+0.04}$ | $4.2{ }_{-1.0}^{+1.1} \times 10^{2}$ | $1.83_{-0.04}^{+0.03}$ |
| NGC0925 | $2.55_{-0.3}^{+0.5} \times 10^{7}$ | $3.5{ }_{-0.10}^{+0.09} \times 10^{9}$ | $6.00_{-0.17}^{+0.15} \times 10^{-1}$ | $0.88_{-0.1}^{+0.1} \times 10^{1}$ | $0.288_{-0.1}^{+0.05}$ | $3.1_{-1.0}^{+2.9} \times 10^{2}$ | $2.50{ }_{-0.11}^{+0.01}$ |
| N | $8.33_{-0.6}^{+0.7} \times 10^{7}$ | $4.22_{-0.11}^{+0.10} \times 10^{10}$ |  | $3.11_{-0.3}^{+0.2} \times 10^{1}$ | $0.72_{-0.1}^{+0.22}$ | $1.5{ }_{-0.3}^{+0.4} \times 10^{6}$ | $2.30_{-0.04}^{0+.11}$ |
| NGC126 | $8.44_{-0.5}^{+0.0} \times 10^{6}$ | $2.8{ }_{-0.09}^{+0.10} \times 10^{10}$ | $4.9{ }_{-0.16}^{+0.16}$ | $21.9_{-2.3}^{+1.7} \times 10^{1}$ | - ${ }_{-0.7}^{+0.63}$ | $6.1_{-0.5}^{+0.8} \times 10^{6}$ | 2 |
| NGC1291 | $1.6_{-0.2}^{+0.3} \times 10^{7}$ | $2.22_{-0.08}^{+0.07} \times 10^{9}$ | $3.7_{-0.14}^{+0.12} \times 10^{-1}$ | $0.88_{-0.1}^{+0.1} \times 10^{1}$ | - $3_{-0.1}^{+0.09}$ | $1.22_{-1.4}^{+2 \text { 24.08.6 }} \times 10$ | $2.50_{-0.28}^{+0.00}$ |
| NGC1 | $9.44_{-0.8}^{+1.2} \times 10^{6}$ | $5.9_{-0.20}^{+0.19}$ | -0.1 | $4.00_{-0.5}^{+0.4} \times 10^{1}$ | $3.55_{-1.1}^{+1.45}$ | 1.0 | $2.45{ }_{-0.52}^{+0.00}$ |
| NGC137 | $1.55_{-0.1}^{+0.2} \times 10^{6}$ | $1.4{ }_{-0.03}^{+0.04} \times$ | 2.5 | $63.88_{-9.0}^{+5.5} \times 10^{1}$ | $5.52_{-1}^{+1}$ | $1.0_{-0.04}^{+0.0 .0} \times 10^{7}$ | $2.04{ }_{-0.02}^{+0.01}$ |
| GC1404 |  |  |  |  |  |  |  |
| IC | $4.0_{-0.3}^{+0.3} \times$ | $1.6{ }^{+0.04}$ | $2.88^{+0.0}$ | 2 | 75 | $4_{-0.4}^{+0.9} \times$ | 00 |
| NGC1482 | $2.5_{-0.2}^{+0.2} \times 10^{7}$ | $5.1_{-0.14}^{+0.13} \times 10^{10}$ | $8.8{ }_{-0.24}^{+0.22}$ | $12.2{ }_{-1.1}^{+1.1} \times 10^{1}$ | 2.65 | $2.2{ }_{-0.5}^{+0.4} \times 10^{6}$ | 2.27 |
| NGC1512 | $3.11_{-0.6}^{+1.1} \times 10^{7}$ | $3.3{ }_{-0.11}^{+0.09} \times 10^{9}$ | $5.7_{-0.18}^{+0.16} \times 10^{-1}$ | $0.7_{-0.2}^{+0.2} \times 10^{1}$ | $0.144_{-0.1}^{+0.07}$ | $3.4{ }_{-2.8}^{+4.0} \times 10^{2}$ | $2.19_{-0.18}^{+0.12}$ |
| NG | $4.88_{-0.3}^{+0.5} \times 10^{7}$ | $1.33_{-0.03}^{+0.03} \times$ | $2.33_{-0.06}^{+0.06}$ | $16.5{ }_{-1.7}^{+1.4} \times 10^{1}$ | $4.49_{-0.7}^{+0.63}$ | $3.6_{-1.9}^{+2.0 .0} \times 10^{5}$ | $2.37_{-0.03}^{-0.03}$ |
| HoII | $7.6_{-0.5}^{+0.7} \times 10^{4}$ | $5.1{ }_{-0.13}^{+0.12} \times 10^{7}$ |  | $4.66_{-0.5}^{+0.7} \times 10^{1}$ | 10 | $4.6{ }_{-0.7}^{+1.9} \times 10^{5}$ | $2.21_{-0.02}^{+0.02}$ |
| DDO05 | $1.0_{-0.1}^{+0.2} \times 10^{4}$ | $1.1{ }_{-0.03}^{+0.03} \times 10^{7}$ | $2.0{ }_{-0.05}^{++0.25} \times 10^{-3}$ | 7.0 ${ }_{-0.9}^{+0.8} \times 10^{1}$ | ${ }^{4}$ | $5.6{ }_{-5.5}^{+0.1} \times 10^{5}$ | $2.30_{-0.12}^{+0.04}$ |
| NGC2798 | $1.22_{-0.1}^{+0.1} \times 10^{7}$ | $4.0_{-0.11}^{+0.11} \times 10^{10}$ | $6.9{ }_{-0.19}$ | $20.4_{-1.8}^{+1.6} \times 10^{1}$ | 3.59 | 6.15 |  |
| NGC2 | $5.9{ }_{-0.6}^{+0.5} \times 10^{7}$ | $9.00_{-0.33}^{+0.32} \times 10^{9}$ | $1.6-0.00$ | $0.9_{-0.1}^{+0.1} \times 10^{1}$ | $0.60{ }^{+0.07}$ | 1.0 |  |
| NGC2 | $3.4{ }_{-0.6}^{40.8}$ | 3.8 | ${ }_{-0.20}^{+0.19} \times$ | $0.7_{-0.2}^{+0.2} \times 10^{1}$ |  | $33^{+1.0} \times 10^{1}$ |  |
| HoI | 2.4 | $1.2{ }_{-0.03}^{+0.03} \times 10^{7}$ | $2.1_{-0.06}^{+0.06} \times$ | $0.3{ }_{-0.03}^{+0.03} \times 10^{1}$ | 0 | $2.6_{-0.4}^{+0.5} \times 10^{2}$ | $1.76_{-0.02}^{+0.02}$ |
| N | $1.9_{-0.1}^{+0.2} \times$ | $7.3{ }_{-0.19}^{+0.21} \times 10^{8}$ | $1.3{ }_{-0.03}^{+0.04} \times 10$ | - | $0.966_{-0.1}^{+0.0 .07}$ | $1.6{ }_{-3.1}^{+999995.3} \times 10^{1}$ | $2.50_{-0.00}^{0+0.00}$ |
| NGC304 | $7.7_{-1.5}^{+1.7} \times 10^{6}$ | $3.6{ }_{-0.07}^{+0.08} \times 10^{9}$ | $6.22_{0.13}^{+0.14} \times 10$ | $2.88_{-0.5}^{+0.8} \times 10^{1}$ | 14 | $4.8{ }_{-2.3}^{+471.5} \times 10^{4}$ | 2.11 |
| NGC307 | $1.11_{-0.1}^{+0.2} \times 10^{6}$ | $7.5{ }_{-0.18}^{+0.17} \times 10^{8}$ | $1.3{ }_{-0.03}^{+0.03} \times 10$ | $3.99_{-0.6}^{+0.6} \times 10^{1}$ | .98 ${ }_{-0.3}^{+0.39}$ | $1.0_{-0.4}^{+0.5} \times 10^{3}$ | 2.2 |
| M81dwB | $2.88_{-0.4}^{+3.1} \times 10^{4}$ | $4.7{ }_{-017}^{+0.11} \times 10^{6}$ | $8.1{ }_{-0.30}^{1+0.19} \times 10^{-1}$ | $1.1_{-0.2}^{+0.0} \times 10^{1}$ | 0 | $1.5{ }_{-1.7}^{5+1.6} \times 10^{4}$ |  |
| NGC3 | $1.8{ }_{-0.2}^{+0.4} \times 10^{7}$ | $6.4{ }_{-0.17}^{+0.16} \times 10^{9}$ | 0.30 | $2.1_{-0.2}^{+0.5} \times 10^{1}$ | 0.80 | $3 .{ }^{-2.5}{ }_{-2.7}^{-9.7} \times 10^{1}$ | .50 |
| NGC318 | $3.22_{-0.3}^{+0.5} \times 10^{7}$ | $8.0{ }_{-0.20}^{+0.17} \times 10^{9}$ |  | $1.5_{-0.2}^{+0.2} \times 10^{1}$ | $1.000_{-0.4}^{+0.23}$ | $1.4{ }_{-0.3}$ | 2.50 |
| NGC3198 | $3.88_{-0.4}^{+0.6} \times 10^{7}$ | $7.4{ }_{-0.17}^{+0.17} \times 10^{9}$ | $1.3_{-0.03}^{+0.03}$ | $1.11_{-0.1}^{+0.2} \times 10^{1}$ | $0.42_{-0.1}^{+0.40}$ | $2.8{ }_{-1.7}^{+8.0} \times 10^{1}$ | 2.50 |
| IC2574 | $2.88_{-0.6}^{+2.0} \times 10^{6}$ | $2.1{ }^{-0.08}$ | $3.6_{-0.15}^{+0.10} \times 10^{-2}$ | $0.5_{-0.2}^{+0.1} \times 10^{1}$ | , | $3.5{ }_{-2.7}^{+1.0} \times 10^{2}$ | 1.9 |
| NGC3265 | $1.88_{-0.2}^{+0.3} \times 10^{6}$ | $2.6{ }_{-0.06}^{+0.066} \times 10^{9}$ | $4.5_{-0.11}^{+0.11} \times 10^{-1}$ | $8.88_{-1.3}^{+0.9} \times 10^{1}$ | $1.455_{-0.2}^{+0.24}$ | $3.0{ }_{-2.3}^{+1.0} \times 10^{6}$ | $2.18{ }_{-0.02}^{+0.02}$ |
| NG | $1.9{ }_{-0.1}^{+0.1} \times 10^{7}$ | $6.9 .9{ }_{-0.18}^{+0.17} \times 10^{9}$ |  | $2.1_{-0.2}^{+0.2} \times 10^{1}$ | $0.700_{-0.1}^{+0.09}$ | $1.4{ }_{-1.2}^{+0.3} \times 10^{6}$ |  |
| NGC3521 | $7.11_{-0.5}^{+1.2} \times 10^{7}$ | $3.1_{-0.08}^{+0.07} \times 10^{10}$ |  | $2.5{ }^{+0.2}+0.2 \times 10^{1}$ | $1.57{ }_{-0.6}^{+0.15}$ | $3.2{ }^{-1.5}$ | 2.50 |
| NGC3621 | $1.99_{-0.2}^{+0.3} \times 10^{7}$ | $7.0{ }_{-0.16}^{+0+17} \times 10^{9}$ | $1.2_{-0.03}^{+0.0 .03}$ | $2.1_{-0.3}^{+0.4} \times 10^{1}$ | $0.89{ }_{-0.2}^{+0.32}$ | $1.44_{-1.0}^{+3.2} \times 10^{1}$ | $2.50_{-0.00}^{+0.00 .00}$ |

Table 7.5. SED parameters: $M_{d u s t}, L_{T I R}, \mathrm{SFR},\langle U\rangle, U_{\text {min }}, \Delta U$ and $\alpha$ for the KINGFISH sample.

| Source | $M_{\text {dust }}$ | $L_{T I R}$ | SFR <br> [M <br>  <br>  <br> $\left[M_{\odot}\right]$ | $\left[L_{\odot}\right]$ |
| :--- | :---: | :---: | :---: | :---: | :---: | :---: | :---: |

Table 7.6. SED parameters: $f_{v s g}, f_{P A H}, f_{i o n}$ and $M_{\star}$ for the KINGFISH sample.

| Source | $f_{v s g}$ | $f_{P A H}$ | $f_{\text {ion }}$ | $\begin{gathered} M_{\star} \\ {\left[M_{\odot}\right]} \end{gathered}$ |
| :---: | :---: | :---: | :---: | :---: |
| NGC0337 | $0.47_{-0.08}^{+0.07}$ | $0.47_{-0.02}^{+0.03}$ | $0.96{ }_{-0.08}^{+0.04}$ | $2.5_{-0.03}^{+0.03} \times 10^{10}$ |
| NGC0584 | - | - | - |  |
| NGC0628 | $0.944_{-0.08}^{+0.06}$ | $1.40_{-0.06}^{+0.06}$ | $0.34_{-0.04}^{+0.04}$ | $2.8{ }_{-0.04}^{+0.03} \times 10^{10}$ |
| NGC0855 | 1.00 | $0.16_{-0.01}^{+0.01}$ | 1.00 | $2.99_{-0.04}^{+0.03} \times 10^{9}$ |
| NGC0925 | $0.96_{-0.06}^{+0.06}$ | $0.78_{-0.05}^{+0.04}$ | $0.44_{-0.06}^{+0.06}$ | $1.7_{-0.02}^{+0.02} \times 10^{10}$ |
| NGC1097 | $0.74_{-0.23}^{+0.42}$ | $0.811_{-0.08}^{+0.05}$ | $0.41_{-0.07}^{+0.09}$ | $1.7_{-0.04}^{+0.03} \times 10^{11}$ |
| NGC1266 | $0.10_{-0.07}^{+0.06}$ | $0.12_{-0.01}^{+0.01}$ | 0.00 | $4.0_{-0.05}^{+0.05} \times 10^{10}$ |
| NGC1291 | $0.96_{-0.28}^{+0.06}$ | 0.00 | 1.00 | $1.6_{-0.03}^{+0.03} \times 10^{11}$ |
| NGC1316 | $1.55_{-0.40}^{+0.14}$ | 0.00 | $1.00_{-0.52}^{+0.48}$ | $7.5_{-0.10}^{+0.11} \times 10^{11}$ |
| NGC1377 | $0.00_{-0.00}^{+0.11}$ | $0.43_{-0.02}^{+0.04}$ | 1.00 | $2.2_{-0.04}^{+0.03} \times 10^{10}$ |
| NGC1404 | - |  |  |  |
| IC0342 | $0.677_{-08}^{+0.10}$ | $1.599_{-005}^{+0.06}$ | 0.00 | $3.6_{-0.05}^{+0.04} \times 10^{10}$ |
| NGC1482 | $0.55_{-0.07}^{+0.07}$ | $0.90_{-0.03}^{+0.03}$ | $0.42_{-0.02}^{+0.03}$ | $4.7_{-0.06}^{+0.05} \times 10^{10}$ |
| NGC1512 | 1.00 | $0.46_{-0.03}^{+0.03}$ | 1.00 | $3.9{ }^{-0.05}+10.040^{-0.04} \times 10$ |
| NGC2146 | $0.93_{-0.05}^{+0.05}$ | $1.04_{-0.03}^{+0.04}$ | $0.56_{-0.02}^{+0.02}$ | $9.5{ }_{-0.11}^{+0.12} \times 10^{10}$ |
| HoII | $0.00_{-0.01}^{+0.01}$ | 0.00 | $0.87_{-0.24}^{+0.13}$ | $6.44_{-0.08}^{+0.09} \times 10^{8}$ |
| DDO053 | $0.93_{-0.11}^{+0.12}$ | $0.13_{-0.01}^{+0.01}$ | 1.00 | $4.3_{-0.05}^{+0.05} \times 10^{7}$ |
| NGC2798 | $0.00_{-0.00}^{+0.001}$ | $0.53_{-0.02}^{+0.02}$ | $0.35_{-0.03}^{+0.03}$ | $4.0_{-0.05}^{+0.05} \times 10^{10}$ |
| NGC2841 | $1.00_{-0.40}^{+0.11}$ | $0.78_{-0.08}^{+0.46}$ | $0.87_{-0.57}^{+0.17}$ | $2.0_{-0.03}^{+0.03} \times 10^{11}$ |
| NGC2915 | $1.44_{-0.09}^{+0.08}$ | $0.35_{-0.03}^{+0.03}$ | 0.00 | $6.1_{-0.07}^{+0.06} \times 10^{8}$ |
| HoI | 1.00 | $0.18_{-0.01}^{+0.01}$ | 1.00 | $1.3{ }^{-0.02} \times 10^{8}$ |
| NGC2976 | $1.31_{-0.41}^{+0.09}$ | $0.88_{-0.06}^{+0.08}$ | $0.49_{-0.09}^{+0.08}$ | $3.5_{-0.05}^{+0.05} \times 10^{9}$ |
| NGC3049 | $0.588_{-0.57}^{+0.10}$ | $0.89_{-0.17}^{+0.04}$ | $0.37_{-0.06}^{+0.07}$ | $9.9{ }_{-0.14}^{+0.16} \times 10^{9}$ |
| NGC3077 | $1.39_{-0.11}^{+0.10}$ | $0.72_{-0.05}^{+0.05}$ | $0.41_{-0.08}^{+0.09}$ | $4.88_{-0.05}^{+0.05} \times 10^{9}$ |
| M81dwB | $0.76_{-0.31}^{+0.83}$ | $0.11_{-0.08}^{+0.03}$ | 1.00 | $5.9_{-0.13}^{+0.08} \times 10^{7}$ |
| NGC3190 | $0.52_{-0.06}^{+0.05}$ | $1.155_{-0.06}^{+0.066}$ | $0.155_{-0.05}^{+0.05}$ | $1.0_{-01}^{+0.01} \times 10^{11}$ |
| NGC3184 | $1.11_{-0.07}^{+0.06}$ | $1.12_{-0.05}^{+0.05}$ | $0.87_{-0.06}^{+0.07}$ | $5.1_{-0.07}^{+0.05} \times 10^{10}$ |
| NGC3198 | $1.400_{-0.10}^{+0.08}$ | $1.11_{-0.06}^{+0.06}$ | $0.32_{-0.05}^{+0.05}$ | $3.88_{-0.06}^{+0.05} \times 10^{10}$ |
| IC2574 | $1.13_{-0.08}^{+0.11}$ | $0.00_{-0.00}^{+0.003}$ | 0.00 | $1.22_{-0.02}^{+0.02} \times 10^{9}$ |
| NGC3265 | $0.45_{-0.19}^{+0.22}$ | $0.67_{-0.04}^{+0.06}$ | $0.46_{-0.07}^{+0.07}$ | $7.1_{-0.09}^{+0.09} \times 10^{9}$ |
| NGC3351 | $1.66_{-0.15}^{+0.30}$ | $0.53_{-0.03}^{+0.04}$ | $1.00_{-0.10}^{+0.00}$ | $5.0_{-0.06}^{+0.06} \times 10^{10}$ |
| NGC3521 | $0.79_{-0.07}^{+0.066}$ | $1.54_{-0.06}^{+0.06}$ | $0.45_{-0.04}^{+0.05}$ | $1.7_{-0.02}^{+0.02} \times 10^{11}$ |
| NGC3621 | $1.00_{-0.09}^{+0.07}$ | $1.54_{-0.07}^{+0.066}$ | $0.32_{-0.04}^{+0.04}$ | $2.88_{-0.04}^{+0.03} \times 10^{10}$ |
| NGC3627 | $0.90_{-0.09}^{+0.10}$ | $1.155_{-0.05}^{+0.04}$ | $0.39_{-0.05}^{+0.04}$ | $1.1{ }_{-0.01}^{+0.01} \times 10^{11}$ |
| NGC3773 | $0.20_{-0.28}^{+0.34}$ | $0.85{ }_{-0.05}^{+0.05}$ | $0.01_{-0.03}^{+0.06}$ | $2.22_{-0.02}^{+0.03} \times 10^{9}$ |
| NGC3938 | $1.00_{-0.08}^{+0.066}$ | $1.23_{-0.06}^{+0.05}$ | $0.34_{-0.05}^{+0.04}$ | $6.5_{-0.09}^{+0.08} \times 10^{10}$ |
| NGC4236 | $0.00_{-0.01}^{+0.01}$ | $0.23_{-0.02}^{+0.02}$ | $1.00_{-0.09}^{+0.00}$ | $4.2_{-0.05}^{+0.05} \times 10^{9}$ |
| NGC4254 | 1.00 | $1.48_{-0.07}^{+0.07}$ | $0.46_{-0.04}^{+0.04}$ | $8.7{ }_{-0.14}^{+0.11} \times 10^{10}$ |
| NGC4321 | $1.00_{-0.08}^{+0.07}$ | $1.23_{-0.06}^{+0.06}$ | $0.43_{-0.05}^{+0.04}$ | $1.3_{-0.02}^{+0.02} \times 10^{11}$ |
| NGC4536 | $0.67_{-0.11}^{+0.088}$ | $0.83{ }_{-0.04}^{+0.07}$ | $0.45_{-0.05}^{+0.05}$ | $5.3_{-0.06}^{+0.06} \times 10^{10}$ |
| NGC4559 | $0.78_{-0.29}^{+0.07}$ | $1.10_{-0.06}^{+0.09}$ | $0.24_{-0.05}^{+0.05}$ | $1.2_{-0.02}^{+0.02} \times 10^{10}$ |
| NGC4569 | $1.03_{-0.10}^{+0.10}$ | $1.511_{-0.05}^{+0.05}$ | 0.00 | $5.8{ }_{-0.07}^{+0.07} \times 10^{10}$ |
| NGC4579 | 1.00 | $0.84_{-0.10}^{+0.10}$ | $0.37_{-0.12}^{+0.17}$ | $1.9{ }_{-0.02}^{+0.02} \times 10^{11}$ |
| NGC4594 | $0.00_{-0.00}^{+0.38}$ | $1.55_{-0.28}^{+0.10}$ | 0.00 | $2.4{ }_{-0.03}^{+0.03} \times 10^{11}$ |
| NGC4625 | $1.24_{-0.09}^{+0.08}$ | $1.40_{-0.07}^{+0.07}$ | $0.40_{-0.05}^{+0.05}$ | $2.99_{-0.04}^{+0.03} \times 10^{9}$ |
| NGC4631 | $0.70_{-0.15}^{+0.25}$ | $1.01_{-0.06}^{+0.06}$ | $0.40_{-0.05}^{+0.07}$ | $3.8{ }_{-0.05}^{+0.05} \times 10^{10}$ |
| NGC4725 | $0.84_{-0.35}^{+0.19}$ | $1.20_{-0.23}^{+0.44}$ | $0.63_{-0.34}^{+0.38}$ | $1.3{ }_{-0.02}^{+0.02} \times 10^{11}$ |
| NGC4736 | $0.89_{-0.13}^{+0.09}$ | $1.32_{-0.07}^{+0.07}$ | $0.18_{-0.05}^{+0.05}$ | $5.6_{-0.07}^{+0.06} \times 10^{10}$ |
| DDO154 | - | - | - | - |
| NGC4826 | $0.19_{-0.53}^{+0.10}$ | $1.12_{-0.08}^{+0.12}$ | $0.04_{-0.06}^{+0.08}$ | $5.6_{-0.07}^{+0.07} \times 10^{10}$ |
| DDO165 | - | - | - | - |

Table 7.6. SED parameters: $f_{v s g}, f_{P A H}, f_{i o n}$ and $M_{\star}$ for the KINGFISH sample.

| Source | $f_{\text {vsg }}$ | $f_{P A H}$ | $f_{\text {ion }}$ | $M_{\star}$ <br> $\left[M_{\odot}\right.$ |
| :--- | :---: | :---: | :---: | :---: |
| NGC5055 | $0.79_{-0.08}^{+0.07}$ | $1.42_{-0.06}^{+0.07}$ | $0.25_{-0.04}^{+0.05}$ | $9.8_{-0.13}^{+0.11} \times 10^{10}$ |
| NGC5398 | $2.67_{-0.10}^{+0.10}$ | $0.65_{-0.04}^{+0.06}$ | 0.00 | $1.1_{-0.01}^{+0.01} \times 10^{9}$ |
| NGC5408 | 1.00 | $0.02_{-0.005}^{+0.01}$ | 1.00 | $1.1_{-0.01}^{+0.01} \times 10^{9}$ |
| NGC5457 | $1.46_{-0.08}^{+0.07}$ | $0.84_{-0.05}^{+0.04}$ | $0.61_{-0.06}^{+0.07}$ | $7.1_{-0.09}^{+0.07} \times 10^{10}$ |
| NGC5474 | $0.72_{-0.06}^{+0.08}$ | $0.95_{-0.04}^{+0.04}$ | 0.00 | $2.4_{-0.02}^{+0.02} \times 10^{9}$ |
| NGC5713 | $0.40_{-0.05}^{+0.05}$ | $0.57_{-0.03}^{+0.03}$ | $0.78_{-0.04}^{+0.05}$ | $5.9_{-0.07}^{+0.07} \times 10^{10}$ |
| NGC5866 | $0.18_{-0.04}^{+0.04}$ | $0.81_{-0.04}^{+0.04}$ | 0.00 | $1.2_{-0.01}^{+0.01} \times 10^{11}$ |
| NGC6946 | $0.99_{-0.09}^{+0.12}$ | $1.24_{-0.06}^{+0.04}$ | $0.45_{-0.04}^{+0.05}$ | $9.3_{-0.13}^{+0.14} \times 10^{10}$ |
| NGC7331 | $1.02_{-0.42}^{+0.12}$ | $1.22_{-0.08}^{+0.09}$ | $0.28_{-0.06}^{+0.07}$ | $2.2_{-0.03}^{+0.03} \times 10^{11}$ |
| NGC7793 | $0.71_{-0.06}^{+0.05}$ | $0.92_{-0.04}^{+0.04}$ | $0.73_{-0.06}^{+0.06}$ | $8.3_{-0.11}^{+0.09} \times 10^{9}$ |

When no errors are given, it means that either this parameter was fixed in the fit or that the parameter reached the minimum or maximum value allowed for this parameter. If fixed: $f_{v s g}=1, f_{P A H}=0.0, f_{i o n}=0.5$, and $f_{v s g} \geq 0.0, f_{P A H} \geq 0.0$ and $f_{\text {ion }} \in[0.0,1.0]$.

## Chapter 8

## Characterisation of the dust properties

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In this Chapter, we analyse the results of the SED model fits that were performed and presented in the previous Chapter. In the first section, we demonstrate the importance of Herschel observations to properly constrain the model solutions for the dust properties. In Section 8.2, we explore the various dust properties that were exploited in Chapter 5 and study their variations with metallicity now using a realistic description of dust. In Section 8.3 we investigate the submm excess, and the conditions for its presence with other galactic properties.

Throughout this Chapter we would like to remind the reader that the results presented here have been obtained assuming the SPIRE beam areas and calibration values from September 2012. These values are currently ${ }^{1}$ being updated by the SPIRE ICC with a new model for the beam. The new numbers for the beam areas may imply a difference for the SPIRE fluxes of the extended sources between 5 and $10 \%$. This will necessarily affect the quantitative values of the results that are presented here, especially for Section 8.3. However, we believe that the qualitative trends presented in Sections 8.1 and 8.2 would not be too affected. Thus we refer any future reader to the papers that should be published soon ${ }^{2}$ for the updated results.

### 8.1 Input from Herschel: Importance of submm observational constraints

### 8.1.1 On the dust mass

Galametz et al. (2011) demonstrated the importance of submm constraints beyond $160 \mu \mathrm{~m}$ on a sample of 52 galaxies and found that the dust mass estimate can be biased in a non-trivial way, apparently depending on the metallicity of the galaxy, if no submm constraints are used in the modelling. Dale et al. (2012) showed with the KINGFISH sample that the dust masses estimated without Herschel constraints were consistent within a factor of 2 with the dust masses estimated with both Herschel and Spitzer constraints. Nonetheless, they noted that for galaxies with $12+\log (\mathrm{O} / \mathrm{H}) \leq 8.0$, dust masses tend to be underestimated if no Herschel constraints are used in the fit, in agreement with the results at low metallicities of Galametz et al. (2011).

We perform the test on both DGS and KINGFISH samples and estimate the dust masses using only Spitzer constraints, $M_{d}(S)$, and using both Herschel and Spitzer constraints, $M_{d}(S+H)$. We restrict our sample to galaxies detected in the MIPS 160 band which reduces the number of lowmetallicity galaxies for this comparison. The results are presented in Fig. 8.1 as a function of metallicity. Additionally we colour code the galaxies depending on their FIR colour, $F_{70} / F_{160}$, that can be indicative of the global dust temperature.

The median ratio of the two dust mass estimates: $M_{d}(S) / M_{d}(S+H)$ is 1.1 for the DGS sample and 1.5 for KINGFISH galaxies with a higher dispersion in the values for the KINGFISH sample (Fig. 8.1): 0.52 dex compared to 0.37 dex for the DGS galaxies. The dust masses estimated without Herschel constraints are overestimated for cold high-metallicity galaxies (i.e., $12+\log (\mathrm{O} / \mathrm{H}) \gtrsim 8.1$ and $F_{70} / F_{160}<1$ ) in agreement with the conclusions of Galametz et al. (2011). We also find that for low-metallicity galaxies the dust masses tend to be lower when estimated without Herschel constraints (for galaxies with $12+\log (\mathrm{O} / \mathrm{H}) \lesssim 8.1$ the median ratio drops to 0.9 ). This is consistent with the trends presented in Galametz et al. (2011) and Dale et al. (2012).

Herschel data provide constraints to the submm slope of the dust SED by probing the cold dust component and thus constrain the minimum starlight intensity parameter, $U_{\text {min }}$. This parameter significantly influences the final dust mass: the lower $U_{\text {min }}$, the higher the dust mass. The middle panel of Fig. 8.1 shows the ratio of the two estimates of $U_{\text {min }}$ and we see that the $U_{\text {min }}$ parameter

[^44]tends to be underestimated for high-metallicity galaxies when no submm constraints are used, thus leading to the observed overestimation of the dust mass. The fit without submm constraints indeed needs to invoke more cold dust in order to reproduce the available constraints. However, note that this is model-dependent: we use a model with a distribution of starlight intensity between $U_{\min }$ and $U_{\text {min }}+\Delta U$, whereas Dale et al. (2012) use the Draine \& Li (2007) model which also includes a single starlight intensity component modelled by a $\delta$-function: $\delta\left(U-U_{\min }\right)$. The results for $U_{\min }$ will necessarily be different, but this will not change the fact that we need submm constraints to accurately determine the dust mass parameter.


Fig. 8.1. Comparison of the dust mass, $M_{d}$, minimum starlight intensity, $U_{m i n}$, and TIR luminosity, $L_{T I R}$, estimated with both Herschel and Spitzer constraints (S+H) and only Spitzer (S). The crosses indicate DGS galaxies and the downward triangles show KINGFISH galaxies. The colours code the FIR colour 70/160 estimated with PACS fluxes.

### 8.1.2 On the TIR luminosity

From the previous section, we see that submm constraints are crucial to properly probe the cold dust component of the SED. However, the cold dust is not the main contributor to the TIR luminosity ${ }^{3}$ and thus we do not expect a big difference between $L_{T I R}(S+H)$ and $L_{T I R}(S)$. Indeed the median ratio for both samples is 1.11 with a dispersion of 0.05 dex. For the DGS sample alone, the median

[^45]ratio is 0.98 , and 1.15 for the KINGFISH sample with, again, a slightly higher dispersion: 0.05 dex compared to 0.04 dex for the DGS galaxies. On the bottom panel of Fig. 8.1 we note that the TIR luminosity for cold galaxies with $F_{70} / F_{160}<1$, is systematically overestimated by $\sim 16 \%$ when the submm constraints are omitted.

For the three parameters we investigate in this section, the sampling of the peak of the SED determines if the parameter needs submm constraints to be accurately determined. Indeed, galaxies with $F_{70} / F_{160}>1$ (blue points in Fig. 8.1) have the peak of their SED at shorter wavelengths, around $70 \mu \mathrm{~m}$, and the last MIPS point at $160 \mu \mathrm{~m}$ already provides a sufficient constraint on the Rayleigh-Jeans slope of the dust SED. On the contrary, if $F_{70} / F_{160}<1$, the last constraint available before the advent of Herschel did not enable a good sampling of the peak and of the Rayleigh Jeans slope of the dust SED and the estimation of the dust parameters was more uncertain. This is illustrated in Fig. 8.2 for two galaxies with $F_{70} / F_{160}>1$ (Haro 2) and $F_{70} / F_{160}<1$ (NGC 6946).


Fig. 8.2. Comparison between the SEDs obtained with and without Herschel constraints for Haro 2 (left), a DGS galaxy with $F_{70} / F_{160}>1$, and NGC6946 (right), a KINGFISH galaxy with $F_{70} / F_{160}<1$. Herschel constraints are indicated with the purple crosses, MIPS with orange stars and other ancillary data with orange crosses. The SED models obtained with and without Herschel constraints are shown with the solid $(\mathrm{S}+\mathrm{H}$ model) and dashed ( S model) lines respectively. Note that for Haro2, the submm point at $850 \mu \mathrm{~m}$ is not taken into account in both fits ( $\mathrm{S}+\mathrm{H}$ and S ) as explained in Section 7.2 (zero weight): the last constraint for the (S) model is the MIPS point at $160 \mu \mathrm{~m}$.

### 8.2 Dust from low-metallicity environments to "normal" galaxies

### 8.2.1 Importance of a realistic description of dust

In Chapter 5 we conducted a first study of the dust properties in low-metallicity environments using a modified blackbody (modBB) model to describe the dust emission. However, in a modBB model the constraints from the dust emitting at short wavelengths, tracing the warmer dust contribution, are not taken into account and this can affect the total dust mass and properties.

Fig. 8.3 (left) shows the ratios of the dust masses estimated with a modified blackbody model, $M_{\text {dust,BB }}$, and with a semi-empirical SED model, $\mathrm{M}_{\text {dust }}$, for the DGS and KINGFISH samples, as
a function of metallicity and colour-coded with $\beta_{o b s}$. The masses are from Tables 5.1, 5.2, 7.3 and 7.5 .


Fig. 8.3. Ratios of the dust masses estimated with a modified blackbody model and with a semi-empirical SED model, $M_{d u s t, B B} / M_{d u s t}$, for the DGS (crosses) and KINGFISH (downward triangles) samples, as a function of the metallicity. The colour codes the modBB emissivity index $\beta_{o b s}$ of the sources (from Chapter 5 , Tables 5.1, 5.2). The median error on the ratio is shown with the dark grey point on the right of the plot. On the left panel, $\beta$ is a free parameter in the modBB fits, whereas it is fixed to 2.0 on the right panel.

We see in Fig. 8.3 (left) that the dust mass estimated from a modBB model is misestimated compared to the mass estimated with a more realistic dust model. The median ratio of $M_{d u s t, B B} / M_{d u s t}$ is 0.46 for the total sample and 0.28 and 0.51 for the DGS and KINGFISH samples alone respectively. Our results for KINGFISH are in agreement with Dale et al. (2012) who report a factor of 1.9 difference between $M_{d u s t}$ and $M_{d u s t, B B}$. We note that the underestimation of the dust masses from a modBB model compared to a full semi-empirical dust SED model increases with decreasing metallicity, and that at a given metallicity $M_{d u s t, B B} / M_{d u s t}$ decreases with $\beta_{o b s}$.

However comparing dust masses from a semi-empirical dust model with a given emissivity to dust masses estimated with a modified blackbody where the emissivity index $\beta$ is left varying is inconsistent, as the two estimates of the dust mass arise from two different optical properties: $\beta=2.0$ for the full dust model and $\beta=\beta_{o b s}$ for the modBB model (Bianchi 2013). Thus we do the same test, using this time modBB dust masses estimated with a modified blackbody model where $\beta$ has been fixed to 2.0. The result is shown on the right plot of Fig. 8.3.

When consistently estimated, we find that the ratio $M_{d u s t, B B} / M_{d u s t}$ is systematically $<1.0$ : i.e., the modified blackbody model systematically underestimates the dust mass compared to a semi-empirical SED model. The median ratios are now: 0.55 for both samples, 0.51 for the DGS and 0.58 for KINGFISH galaxies. The trend with metallicity noted on the left panel of Fig.8.3 is not visible anymore on the right panel of Fig. 8.3. This is because for a given galaxy with $\beta_{o b s}<2.0$ (resp. $>2.0$ ), fixing $\beta=2.0$ imposes less (resp. more) emissive grains which translates into a higher (resp. lower) dust mass to account for the same luminosity. Thus galaxies with $\beta_{o b s}<2.0$, mostly metal-poor galaxies, have $M_{d u s t, B B}(\beta=2.0)$ higher than their original modBB dust mass, resulting in an increased $M_{d u s t, B B} / M_{d u s t}$ ratio. On the contrary, galaxies with a $\beta_{o b s}>2.0$, which are mostly metal-rich galaxies, have a $M_{d u s t, B B}(\beta=2.0)$ lower than their original modBB dust mass, resulting in a decreased $M_{d u s t, B B} / M_{d u s t}$ ratio. The observed trend with metallicity in the left panel of Fig. 8.3 is thus only a side effect of the inconsistency of the two dust mass estimates.

Although less obvious than in the left panel of Fig. 8.3, the trend of decreasing $M_{d u s t, B B} / M_{d u s t}$ with decreasing $\beta_{o b s}$ at a given metallicity can still be noted on the right panel of Fig. 8.3. Galaxies
well represented by a low $\beta_{\text {obs }}$ have a shallow Rayleigh Jeans slope of the SED, representative of a broad SED peak. In these cases, imposing $\beta=2.0$ imposes a steeper slope, not representative of the real shape of the SED and thus underestimates even more the dust mass. This shows the importance of the assumption of the dust temperature distribution and its impact on the dust mass: the broader the distribution of temperature, the lower the emissivity index found by a modBB and the larger the dust mass missed by modBB fits.

### 8.2.2 A characteristic temperature distribution

We saw in Chapter 5 that dwarf galaxies present overally warmer dust and potentially a broad dust SED peak. A temperature distribution is thus necessary to properly describe the dust in lowmetallicity galaxies. The dust temperature distribution is directly linked to the range of starlight intensities, $\Delta U$, to which the dust is exposed (see Eq. 1.30 from Chapter 1 linking $U$ and $T_{d}$ ).

Fig. 8.4 shows the range of starlight intensities to which the dust is exposed $\Delta U$ as a function of the average starlight intensity $\langle U\rangle$ for both samples. We use:

$$
\begin{equation*}
T_{d}[K]=17.5 \times U^{1 /(4+\beta)}=17.5 \times U^{1 / 6} \tag{8.1}
\end{equation*}
$$

adapted from Eq. 1.30 to relate the starlight intensity and the dust equilibrium temperature.


Fig. 8.4. $\Delta U$ as a function of the average starlight intensity $\langle U\rangle$ for the DGS (crosses) and KINGFISH (downward triangles) samples. The average dust equilibrium temperature is shown on the top axis. The colour codes the metallicity of the sources. Grey points are where the galaxy should be if no additional MIR modified blackbody is included in the fits.

The median $\langle U\rangle$ is 17.9 for the DGS and 2.2 for the KINGFISH sample. The median $\Delta U$ is $7.7 \times 10^{4}$ for the DGS and $3.5 \times 10^{2}$ for the KINGFISH sample. We see that DGS galaxies have a higher mean starlight intensity and are described by a larger range of starlight intensity values. This manifests in a broader SED peak that is shifted to shorter wavelengths for the DGS galaxies. On the contrary, the more metal-rich environments probed by KINGFISH have lower $\langle U\rangle$ and lower
$\Delta U$. This is consistent with low-metallicity galaxies harbouring warmer dust than more-metal rich environments: in the DGS the median dust temperature estimated with Eq. 8.1 is 28 K , compared to 20 K for the KINGFISH galaxies. Dust grains are not only warmer in dwarf galaxies but they also span a broader range in equilibrium temperatures. This can be estimated from $\Delta U$ and Eq. 8.1: the range of temperatures spanned by the DGS galaxies is $\sim 94 \mathrm{~K}$ while it is 36 K for the KINGFISH galaxies.

The two DGS galaxies with very high $\langle U\rangle$ and low $\Delta U$ are IZw18 ( $\langle U\rangle=390$ and $\Delta U$ $=21)$ and SBS0335-052 $(<U\rangle=1100$ and $\Delta U=1)$. These galaxies indeed peak at very short wavelengths (around $40 \mu \mathrm{~m}$ for IZw18 and $20 \mu \mathrm{~m}$ for SBS0335-052). IZw18 has a rather narrow peak of its dust SED, consistent with its low $\Delta U$, in contrast to SBS0335-052 (see Fig. 7.23). This is due to the additional modified blackbodies that were added in the MIR to obtain a satisfactory fit of the IRS spectrum shape for SBS0335-052. This additional MIR modBB has a great impact on the $\Delta U$ parameter preventing it from reaching higher values. Thus for all DGS galaxies for which we added the MIR modBB, we show with smaller grey points in Fig. 8.4 the place the galaxy would have if we did not add the extra MIR modBB. For most galaxies for which we added the MIR modBB we get a set of parameters that is consistent with the general trend presented for the rest of the sample. For IZw18, the fit without the MIR modBB gives an even higher $\langle U\rangle(\sim 520)$ but a $\Delta U=184$ still lower than that found for the dwarf sample: IZw18 harbours very hot dust spanning a small range of equilibrium temperatures. For SBS0335-052, the new position reflects better the observed SED shape: hot dust $(\langle U\rangle=3340)$ and very broad SED peak $\left(\Delta U=1.6 \times 10^{6}\right)$, and is consistent with the general trend observed for dwarf galaxies.

### 8.2.3 Dust-to-stellar mass ratios

In Chapter 7, we saw that the dust mass is lower in low-metallicity galaxies (Fig. 7.22): there is a difference of about two orders of magnitude between the DGS and KINGFISH median dust masses. We noted the same trend in Chapter 5 when we only used modBB dust masses. In Chapter 5 it also appeared that the proportion of dust mass relative to stellar mass slightly increases between the metal-poor and metal-rich galaxies. However, as shown in Section 8.2.1, the dust masses estimated with a modBB model using a free $\beta$, as we did in Chapter 5, are systematically misestimations of the real dust masses at low-metallicities. Moreover, when estimated with a $\beta=2.0$ modBB, we saw in Chapter 5 that the correlation between the dust-to-stellar mass ratios with metallicity almost vanishes.

Fig. 8.5 shows the dust-to-stellar mass ratio for the DGS and KINGFISH galaxies as a function of metallicity, this time using dust masses from our semi-empirical SED model. The Spearman rank coefficient is 0.06 , confirming what we saw with $\beta=2.0 \operatorname{modBB}$ dust masses. The median dust-to-stellar mass ratios are now very similar for the DGS and KF samples: $0.07 \%$ vs $0.09 \%$. The scatter in the DGS dust-to-stellar mass ratios is 0.6 dex versus 0.4 dex for the KINGFISH galaxies. So the trend observed in Section 5.3 was due to the adopted dust model. Thus on average, lowmetallicity dwarf galaxies contain as much dust as their metal-rich counterparts when compared to their respective stellar masses.

### 8.2.4 $L_{T I R}$-to- $M_{\text {dust }}$ ratios

In Chapter 5 we also saw that the dwarf galaxies emit more in the FIR/submm than more metal-rich galaxies, per unit dust mass ( $\sim 6$ times more for the DGS) by looking at the $L_{F I R} / M_{d u s t, B B}$ ratio. Here again we want to check if this trend still holds when using dust masses estimated from our semi-empirical dust model. Thus we now consider the $L_{\text {FIR }} / M_{\text {dust }}$ ratio, plotted on the left panel of Fig. 8.6.


Fig. 8.5. Dust-to-stellar mass ratios for the DGS (purple crosses) and KINGFISH (orange downward triangles) samples as a function of metallicity. The distribution of $M_{d u s t} / M_{\star}$ is indicated on the side for both samples: plain purple line for DGS and dashed orange line for KINGFISH.

The median $L_{F I R} / M_{\text {dust }}$ is $1.3 \times 10^{3} L_{\odot} / M_{\odot}$ for the DGS sample, with a standard deviation of 0.5 dex, and $2.7 \times 10^{2} L_{\odot} / M_{\odot}$ for the KINGFISH galaxies, with a standard deviation of 0.4 dex. The correlation is a bit weaker than in Chapter 5: we have a Spearman rank coefficient $\rho=-0.49$, but the difference between the DGS and KINGFISH galaxies remains: the DGS galaxies emit $\sim 4.7$ times more in the FIR/submm domain than the KINGFISH sample. Thus this time, only part of the trend observed in Chapter 5 was due to the adopted dust modelling.

The comparison here is performed with $\mathrm{L}_{F I R}$ estimated between 50 and $650 \mu \mathrm{~m}$. Because the peak of the dust SED for the dwarf galaxies is typically around $40-70 \mu \mathrm{~m}$, the estimation of $\mathrm{L}_{\text {FIR }}$ does not consider the whole peak of the dust SED contrary to more-metal rich environments that have a SED peak around $100 \mu \mathrm{~m}$. Thus we expect the $\mathrm{L}_{\text {FIR }} / L_{T I R}$ ratio to be lower for galaxies harbouring warmer dust, i.e., low-metallicity galaxies with a high mean starlight intensity, compared to galaxies with colder dust, i.e., more metal-rich galaxies with low $\langle U\rangle$. This is indeed what we observe on the right panel of Fig. 8.6. The median $L_{F I R} / L_{T I R}$ is 0.5 for the DGS sample, with a standard deviation of 0.14 dex, and 0.7 for the KINGFISH galaxies, with a standard deviation of 0.06 dex. There is indeed a good correlation between $L_{F I R} / L_{T I R}$ and metallicity: $\rho=0.54$, and an even stronger correlation between $L_{F I R} / L_{T I R}$ and $\langle U\rangle: \rho=-0.82$.

Comparing $L_{F I R}$ to the total dust mass is thus not really appropriate as we compare the total dust mass with only a fraction of the power emitted by the dust grains. A more meaningful comparison is presented in Fig. 8.7 where the total IR luminosity, estimated between 1 and 1000 $\mu \mathrm{m}$, is used.

The median $L_{T I R} / M_{\text {dust }}$ is $2.7 \times 10^{3} L_{\odot} / M_{\odot}$ for the DGS sample, with a standard deviation of 0.6 dex , and $3.8 \times 10^{2} L_{\odot} / M_{\odot}$ for the KINGFISH galaxies, with a standard deviation of 0.4 dex. The Spearman rank coefficient for this trend is $\rho=-0.52$, similar to the Spearman rank coefficient obtained for the $L_{\text {FIR }} / M_{\text {dust }}$ relation with metallicity. However the difference between the DGS and KINGFISH galaxies is more important: the DGS emit $\sim 7$ times more in the infrared domain than


Fig. 8.6. (left) $L_{F I R} / M_{d u s t}$ ratios for the DGS (crosses) and KINGFISH (downward triangles) samples as a function of metallicity. The distribution of $L_{F I R} / M_{d u s t}$ is indicated on the side for both samples. (right) $L_{\text {FIR }} / L_{T I R}$ for the DGS (purple crosses) and KINGFISH (orange downward triangles) samples as a function of metallicity. The distribution of $L_{F I R} / L_{T I R}$ is indicated on the side for both samples. For both panels, the colours code the mean starlight intensity of the galaxies.


Fig. 8.7. $L_{T I R} / M_{\text {dust }}$ ratios for the DGS (crosses) and KINGFISH (downward triangles) samples as a function of metallicity. The distribution of $L_{T I R} / M_{\text {dust }}$ is indicated on the side for both samples. The colours code the mean starlight intensity, $\langle U\rangle$, of the galaxies.
the KINGFISH sample per unit dust mass. This is directly linked to the mean starlight intensity $\langle U\rangle$, as shown by the colours on Fig 8.7, where galaxies with a high $\langle U\rangle$ emit more per unit dust mass than galaxies with a lower $\langle U\rangle$, simply reflecting the fact that warmer dust is more luminous. Thus the trend observed in Chapter 5, Section 5.3.4, in the FIR is confirmed and enhanced when we consider the total IR range.

### 8.2.5 Specific star-formation rates

We finalize our view of the dust properties in galaxies by looking at the specific star-formation rate, $S S F R=S F R / M_{\text {star }}$, for our two samples. The SSFR can be seen as a measure of the efficiency of the star formation. The SFR have been estimated from the TIR luminosities in Chapter 7 and presented in Tables 7.3 and 7.5. The result is shown in Fig. 8.8. The median SSFR is $2.7 \times$ $10^{-10} \mathrm{yr}^{-1}$ for the DGS sample with a standard deviation of 0.4 dex , and $5.7 \times 10^{-11} \mathrm{yr}^{-1}$ for the KINGFISH galaxies with a standard deviation of 0.5 dex. There is a small anti-correlation between the SSFR and metallicities: $\rho=-0.35$. Low-metallicity galaxies are thus more efficient in forming stars than more metal-rich galaxies: we have a difference of a factor of $\sim 5$ between our two samples.


Fig. 8.8. Specific star-formation rates for the DGS (purple crosses) and KINGFISH (orange downward triangles) samples as a function of metallicity. The distribution of $S S F R$ is indicated on the side for both samples.

Note that this relation has been obtained using a $L_{T I R}$-calibrated SFR. Other SFR tracers, such as a combination of FUV $+24 \mu \mathrm{~m}$ or $\mathrm{H} \alpha+24 \mu \mathrm{~m}$, could be more appropriate to estimate the SFR in low-metallicity galaxies where the ratio of obscured versus unobscured star formation is lower than in more metal-rich galaxies (DeLooze et al., in prep). Additionally, we remind the reader here that our sample of dwarf galaxies contain only gas-rich star-forming dwarf galaxies and that we do not consider gas-poor dwarf galaxies in this analysis. Gas-poor dwarfs may present a lower SSFR than that observed in our sample and thus the trend presented here may be weaker if gas-poor dwarfs would be included.

### 8.3 The intriguing submm excess

A submm excess has been observed in the past in several low-metallicity galaxies (Galliano et al. 2003, 2005; Dumke et al. 2004; Bendo et al. 2006; Zhu et al. 2009; Galametz et al. 2009; Bot et al. 2010; Grossi et al. 2010; Galametz et al. 2011). It has been called "excess" because the current available models are unable to fully explain the submm emission of these galaxies (see Section 2.2.2). This submm excess has been one of the main sources of uncertainty in the dust modelling of dwarf galaxies for the past few years.

### 8.3.1 Identifying the galaxies with submm excess

For this study we select galaxies that are detected at $500 \mu \mathrm{~m}$ which gives us 78 galaxies out of the initial sample of 109 galaxies. Following the same procedure as in Section 5.4 we compute the residuals at each observed submm wavelength for the DGS and KINGFISH galaxies: $R(500)$, $R(850), R(870)$ and $R(1200)$. For each residual we compute the error on the residual $\Delta R(\lambda)$ with the same method as in Section 5.4 (using the 300 random realisations of the SEDs). If the residual, $R(\lambda)$, is greater than the error on the observed flux density and greater than $\Delta R(\lambda)$, then the galaxy possesses an excess at the wavelength $\lambda$. We define "signal-to-noise ratio" or "intensity" of the excess, $S / N(\lambda)$, the quantity $R(\lambda) / \Delta R(\lambda)$. The results are shown in Table 8.1. The error shown for the residuals is the larger value between the error on the datapoint and $\Delta R(\lambda)$. As the observations at $850 \mu \mathrm{~m}$ and $870 \mu \mathrm{~m}$ are complementary, we regroup them under $R(850-870)$.

Out of the 78 galaxies detected at $500 \mu \mathrm{~m}, 27$ of them (35\%) show an excess at $500 \mu \mathrm{~m}$. Kirkpatrick et al. (2013) investigate the presence of a submm excess at $500 \mu \mathrm{~m}$ for a subsample of the KINGFISH galaxies, including spiral and dwarf/irregular galaxies, using a two modified blackbodies model. They conclude that even if a small excess is present at $500 \mu \mathrm{~m}$ it is compensated by the same amount of $500 \mu \mathrm{~m}$ "deficiencies". In our study, we find that 23 KINGFISH galaxies have an excess with $\mathrm{S} / \mathrm{N}>1.0$, among which 6 with $\mathrm{S} / \mathrm{N}>3.0$; and only 3 galaxies with $\mathrm{S} / \mathrm{N}(500)<-1$, NGC2146, NGC4254 and NGC4569, which are not in the subsample considered by Kirkpatrick et al. (2013). Thus we cannot say in our study that our $500 \mu \mathrm{~m}$ excesses are compensated by "deficiencies". Kirkpatrick et al. (2013) use a two modified blackbodies model which only allows for two dust temperatures with a warm dust component and a cold dust component. We saw in Section 8.2.2, that a temperature distribution is essential to properly model the dust emission, especially in dwarf galaxies, to account for a possibly broad SED peak. We believe that the difference in the adopted dust SED models may be the main reason for their different findings.

32 of these 78 galaxies have been observed at 850 or $870 \mu \mathrm{~m}^{4}$. However, Draine et al. (2007) advised caution when working with SCUBA data for some of the KINGFISH galaxies: galaxies that have been scan-mapped and for which the data processing can remove some diffuse extended emission; galaxies that have been incompletely mapped with SCUBA and one galaxy, NGC4594 where the $850 \mu$ m flux density is dominated by AGN emission. For galaxies where the SCUBA map does not cover the extent of the whole galaxy, Dale et al. (2005) applied a correction factor to account for the missing emission based on the $160 \mu \mathrm{~m}$ emission. But the $160 / 850 \mu \mathrm{~m}$ ratio can vary spatially within the galaxy and Draine et al. (2007) discarded any galaxy with a correction factor greater than 1.6. If we consider all of these criteria, we must remove from our SCUBA sample the following seven galaxies: NGC1097 (corr > 1.6), NGC4254 (scan-mapped and corr > 1.6), NGC4321 (corr $>1.6$ ), NGC4579 (scan-mapped), NGC4594 (dominated by AGN emission), NGC4736 (corr > 1.6) and NGC6946 (scan-mapped). These galaxies have indeed very low SCUBA flux densities compared to the rest of the FIR/submm photometry, with $R(850) / \Delta R(850) \sim-1.0$. Three additional galaxies also have an even lower $R(850) / \Delta R(850)(\leq-1.3)$ : NGC1482, NGC4536 and NGC4569. For these

[^46]three galaxies, the SCUBA flux density may also be underestimated. Thus in the following, we discard the SCUBA residuals for these ten galaxies.

Finally 22 galaxies out of the 78 galaxies detected at $500 \mu \mathrm{~m}$ have reliable 850 or $870 \mu \mathrm{~m}$ observations ( $28 \%$ ) and 11 show an excess at 850 or $870 \mu \mathrm{~m}$ ( $50 \%$ of the 22 galaxies). Among these 11 galaxies with excess at 850 or $870 \mu \mathrm{~m}$, only 3 already have an excess at $500 \mu \mathrm{~m}$ Haro11, NGC0337 and NGC4631. Thus if we consider only galaxies with both observations at 500 and $850 / 870 \mu \mathrm{~m}$ ( 22 galaxies), the proportion of submm excess increases with wavelength from $14 \%$ to $50 \%$.


Fig. 8.9. Distribution of the residuals at $500 \mu \mathrm{~m}$ (top left), 850 and $870 \mu \mathrm{~m}$ (top right), and $1200 \mu \mathrm{~m}$ (bottom), for the DGS and KINGFISH galaxies. The KINGFISH distribution is shown on top of the DGS distribution. For all of the panels, the colours code the $\mathrm{S} / \mathrm{N}$ of the excess: $R(\lambda) / \Delta R(\lambda)$ from -1.6 to 29 . A grey cell means a $\mathrm{S} / \mathrm{N}<1$ and thus no excess.

For the 5 galaxies observed with MAMBO, three galaxies have an excess: IIZw40 and NGC1569 which also show an excess at $850 \mu \mathrm{~m}$ but not at $500 \mu \mathrm{~m}$; and NGC4214 which does not have an excess at $500 \mu \mathrm{~m}$ and for which no reliable measurements are available at 850 or $870 \mu \mathrm{~m}$. The other two are He2-10 which does not show any excess in the submm wavelengths and NGC1140 which is not detected with MAMBO.

The residual distribution at each wavelength is presented in Fig. 8.9. In total, out of the 78 galaxies detected at $500 \mu \mathrm{~m}, 36$ galaxies ( $46 \%$ ) have an excess in the submm for at least one of the considered wavelengths: $500,850 / 870$ and $1200 \mu \mathrm{~m}$.

### 8.3.2 Excess in the submm with a modified blackbody

In this Section, we compare these findings on the $500 \mu \mathrm{~m}$ submm excess determined from a full dust SED model to that obtained when we looked at the $500 \mu \mathrm{~m}$ submm excess using a modBB model with a fixed $\beta=2.0$ (Section 5.4). A $\beta=2.0$ modBB detected 33 submm excess in the 78 galaxies detected at $500 \mu \mathrm{~m}$. We find that for 8 galaxies the $\beta=2.0$ modBB detects an excess which is not present in the full dust SED model, and for two galaxies, the $\beta=2.0$ modBB does not detect an excess whereas the full dust SED model does. However, the modBB fits were performed including the $500 \mu \mathrm{~m}$ point whereas in the full dust SED model, we did not include the $500 \mu \mathrm{~m}$ point in the fit. We perform the test by removing the $500 \mu \mathrm{~m}$ point in the modBB fits. We find that all of the galaxies presenting an excess when using the $500 \mu \mathrm{~m}$ point also present an excess when the $500 \mu \mathrm{~m}$ point is not included, except one: NGC4594, for which the result is now coherent with the full dust SED model (i.e., no excess).

The two galaxies where the submm excess appears with the full dust SED model and not with the $\beta=2.0$ modBB are NGC 3184 and NGC 4725. For NGC3184, the excess is barely detected at $500 \mu \mathrm{~m}: R(500)_{S E D}=8.3 \pm 8.1$, and thus is not significant. For the second galaxy, $R(500)_{B B}=$ $9.0 \pm 10.6$ but $R(500)_{\text {SED }}=12.4 \pm 8.0$. The excess is indeed significant here in the full dust SED model and has been missed by the modBB.

The $\beta=2.0 \operatorname{modBB}$ thus overpredicts the number of excess galaxies by $\sim 20 \%$. On the contrary, a galaxy showing an excess with a full dust SED model also shows an excess with a $\beta=2.0 \operatorname{modBB}$ in $96 \%$ of the cases considered here.


Fig. 8.10. $R(500)_{B B} / \Delta R(500)_{B B}$ as a function of $\beta_{o b s}$ for the 78 galaxies of the DGS (crosses) and KINGFISH samples (downward triangles) detected at $500 \mu \mathrm{~m}$. The colours code if the excess is detected by the $\beta=2.0 \operatorname{modBB}$ only (blue), by the full dust model only (red) or by both (purple). Galaxies with no excess are shown in black. The $\beta_{o b s}=1.4$ and $\beta_{o b s}=2.0$ lines are added to guide the eye. The $R(500)_{B B} / \Delta R(500)_{B B}=1$ threshold is indicated with the dashed line.

Fig. 8.10 shows the strength of the modBB submm excess, $R(500)_{B B} / \Delta R(500)_{B B}$ as a function of the emissivity index, $\beta_{o b s}$, determined in Chapter 5 . We see that all of the galaxies where only the $\beta=2.0$ modBB predicts an excess have $\beta_{o b s} \leq 1.4$. All of the galaxies with $\beta_{o b s} \leq 1.4$ show an excess either only with the $\beta=2.0$ modBB or with both models, except one. This galaxy is M81dwB which is only detected at the $1.8 \sigma$ level at $500 \mu \mathrm{~m}$ and thus for which it would have been difficult to detect any excess. For galaxies with $\beta_{o b s} \leq 1.4$, the excesses detected only with a $\beta=2.0$ modBB (blue points in Fig. 8.10) have average excess signal-to-noise ratios $R(500)_{B B} / \Delta R(500)_{B B}$ $\sim 1.5$, whereas the excesses detected with both models (purple points in Fig. 8.10) are stronger: $R(500)_{B B} / \Delta R(500)_{B B} \sim 2.4$.

A low $\beta_{\text {obs }}$ means a flatter submm Rayleigh Jeans slope of the modBB, indicative of a relatively broad peak of the dust SED, compared to a $\beta=2.0 \operatorname{modBB}$. Our small excess at $500 \mu \mathrm{~m}$ may just be due to the fact that the $\beta=2.0 \operatorname{modBB}$ cannot account for the particular shape of the SED with a broad peak, which is possible on the other hand with the full dust model. That is why our $\beta=2.0$ modBB model detects an excess in some galaxies even if they do not show an excess when modelled with the full dust model.

To determine if one galaxy has a chance of harbouring a submm excess with a full dust model using only FIR/submm measurements, the few steps to follow would be:

- First, fit a modBB (including the $500 \mu \mathrm{~m}$ point) with a free emissivity index to determine $\beta_{o b s}$.
- Then if the obtained $\beta_{o b s}$ is $\lesssim 1.4$, the second step would be to fit a modBB this time with a fixed emissivity index to 2.0 (still including the $500 \mu \mathrm{~m}$ point), obtain the $500 \mu \mathrm{~m}$ residual, $R(500)_{B B}$, and carefully estimate the error on this residual, $\Delta R(500)_{B B}$.
- If the ratio $R(500)_{B B} / \Delta R(500)_{B B}$ is $\gtrsim 2.0$, then there is a very strong chance that this galaxy presents a submm excess when modelled with a full dust model.
- If the obtained $\beta_{o b s}$ is between 1.4 and 2.0 and if the $\beta=2.0$ modBB detects an excess $\left(R(500)_{B B} / \Delta R(500)_{B B}>1.0\right)$, then the galaxy will also present a submm excess when modelled with a full dust model.
- If $\beta_{\text {obs }}$ is greater than 2.0 , then there is a strong chance that the galaxy does not harbour a submm excess when modelled with a full dust model.


### 8.3.3 Characterising the submm excess

## Evolution with wavelength

From Fig 8.11, we see that the submm excess becomes more pronounced as we go to longer wavelengths, from $500 \mu \mathrm{~m}$ to $850 / 870 \mu \mathrm{~m}$ (except for one galaxy: NGC0337). The difference between the $\mathrm{S} / \mathrm{N}$ of the excess at $850 / 870 \mu \mathrm{~m}$ and the $\mathrm{S} / \mathrm{N}$ of the excess at $500 \mu \mathrm{~m}$ is $\sim 6.3$ for galaxies with an excess at $850 / 870 \mu \mathrm{~m}$. For galaxies with excess detected at both wavelengths, the $\mathrm{S} / \mathrm{N}$ of the excess increases by a factor of $\sim 1.5$.

Only two galaxies have both measurements at 850 and $1200 \mu \mathrm{~m}$ and present excess at these wavelengths. The excess in these three galaxies shows three different behaviours: for NGC1569 the excess increases at 1200 (the S/N goes from 1.7 at $850 \mu \mathrm{~m}$ to 2.9 at $1200 \mu \mathrm{~m}$ ), and for IIZw40, the excess peaks at $850 \mu \mathrm{~m}$ (the $\mathrm{S} / \mathrm{N}$ decreases from 11.4 at $850 \mu \mathrm{~m}$ to 5.5 at $1200 \mu \mathrm{~m}$ ). The evolution of the intensity of the excess at wavelengths greater than $\sim 1000 \mu \mathrm{~m}$ appears to be dependant on the galaxy. More measurements at both wavelengths are needed to observe a more general trend.

From the colours in Fig.8.11, we hint that the excess may appear preferentially in lowermetallicity galaxies. Thus we now look at the $\mathrm{S} / \mathrm{N}$ of the excess as a function of metallicity.


Fig. 8.11. Evolution of the $\mathrm{S} / \mathrm{N}$ of the excess from 500 to $1200 \mu \mathrm{~m}$, for the DGS (top) and the KINGFISH (bottom) samples. The stars indicate an excess i.e., $R(\lambda) / \Delta R(\lambda) \geq 1.0$. The dashed lines connect the datapoints for each galaxy if more than one submm observation is available. The dotted line indicates the $R(\lambda) / \Delta R(\lambda)=1.0$ threshold. The galaxies are colour coded by metallicity.

## Evolution with metallicity



Fig. 8.12. Left: (bottom) $R(500) / \Delta R(500)$ as a function of metallicity for the 78 galaxies detected at $500 \mu \mathrm{~m}$. The dotted line indicates the $R(500) / \Delta R(500)=1.0$ threshold. The colours indicate the status of the excess at longer wavelength: no data available (red), data available but no excess (blue), data available and excess (green). On the top panel, we show the distribution of metallicity for the $78500 \mu$ m-detected galaxies in grey and for the galaxies with an excess at $500 \mu \mathrm{~m}$ in black. Right: Same at $850 / 870 \mu \mathrm{~m}$ for galaxies with data at these wavelengths. The symbols now additionally code if the excess was already present at $500 \mu \mathrm{~m}$ (filled symbols) or not (empty symbols).

Fig. 8.12 shows the $\mathrm{S} / \mathrm{N}$ of the excess as a function of metallicity at 500 and $850-870 \mu \mathrm{~m}$. From Fig. 8.12 left, no trend between the $\mathrm{S} / \mathrm{N}$ of the excess at $500 \mu \mathrm{~m}$ and metallicity can be observed: the Spearman rank coefficient is $\rho=-0.08$ : i.e., the intensity of the excess does not depend on metallicity. However, if we consider only the presence versus the absence of submm excess, we see that the submm excess appears preferentially in metal-poor galaxies: $56 \%$ of the excess galaxies have $Z<0.4 Z_{\odot}(12+\log (\mathrm{O} / \mathrm{H})<8.3)$, whereas, in the total distribution of galaxies detected at $500 \mu \mathrm{~m}$, (grey line on Fig. 8.12 (left)), only $46 \%$ of the total number of galaxies are galaxies with $Z<0.4 Z_{\odot}$. Moreover, the proportion of excess galaxies in the [7.5-8.3] metallicity range is $\sim$ $39 \%$ versus $\sim 31 \%$ in the $] 8.3-8.8]$ range. This confirms the first results obtained with the $\beta=2.0$ $\operatorname{modBB}$. However with the $\beta=2.0 \operatorname{modBB}$ we found that the strongest excesses seems to appear in the most metal-poor galaxies, which is not what we see here. This is due to the fact that the $\mathrm{S} / \mathrm{N}$ of the excess for the $\beta=2.0$ modBB model is anti-correlated with $\beta_{o b s}$ (see Fig. 8.10) and that the lowest $\beta_{o b s}$ were observed in the most metal-poor galaxies.

Additionally, if we compare the appearance of the excess at longer wavelength (i.e., green vs blue crosses in Fig.8.12), we see that this trend is confirmed: for galaxies without excess at $500 \mu \mathrm{~m}$, the submm excess appears at longer wavelengths preferentially in low-metallicity galaxies. This is also confirmed by the right plot of Fig. 8.12: $81 \%$ of the excess galaxies at $850(870) \mu \mathrm{m}$ are galaxies
with $Z<0.4 Z_{\odot}(12+\log (\mathrm{O} / \mathrm{H})<8.3)$, whereas, in the total distribution of galaxies detected at $850(870) \mu \mathrm{m}$, (grey line on Fig. 8.12 (right)), only $38 \%$ of the total number of galaxies are galaxies with $Z<0.4 Z_{\odot}$. Moreover, the proportion of excess galaxies in the [8.0-8.3] metallicity range is $\sim 75 \%$ versus $\sim 10 \%$ in the $] 8.3-8.6]$ range. However, here again, no trend between the strength of the excess and metallicity can be noted: the Spearman rank coefficient is $\rho=-0.09$.


Fig. 8.13. $<U>$ as a function of metallicity for the 78 galaxies detected at $500 \mu \mathrm{~m}$. The colours and symbols are the same as in Fig. 8.12.

We see in Fig. 8.12 that there are six galaxies for which $12+\log (\mathrm{O} / \mathrm{H})$ is lower than 8.0 but for which no excess is detected at $500 \mu \mathrm{~m}$. From Fig.8.13, we see that all of these galaxies with $12+\log (\mathrm{O} / \mathrm{H})<8.0$ and that do not harbour an excess at $500 \mu \mathrm{~m}$ have very low mean starlight intensities, i.e., relatively cold dust given their metallicity. Only one galaxy with $12+\log (\mathrm{O} / \mathrm{H})<8.0$ and low $\langle U\rangle$ has an excess: NGC 2915. For this galaxy, we have $R(500) / \Delta R(500)=1.01$ and therefore this excess is not really significant. Thus, from Fig. 8.13, we infer that the value of $\langle U\rangle$, which is related to the average temperature of the dust, discriminates between galaxies with and without submm excess at $500 \mu \mathrm{~m}$ when $12+\log (\mathrm{O} / \mathrm{H}) \leq 8.0$.

## Colour-colour diagrams

We now explore several FIR/submm colour diagrams in order to see if the submm excess galaxies fall on specific regions in these diagrams. Fig. 8.14 shows two FIR/submm colour-colour diagrams: $F_{160} / F_{250}$ as a function of $F_{250} / F_{500}$ on the left column and $F_{160} / F_{250}$ as a function of $F_{500} / F_{850(870)}$ on the right column. On the top left panel of Fig. 8.14, we also plot the median ratio $\frac{F_{160} / F_{250}}{F_{250} / F_{500}}$ which is $\sim 0.37$. We see that all of the galaxies with an excess at $500 \mu \mathrm{~m}$ fall above this line i.e., they have larger $\frac{F_{160} / F_{250}}{F_{250} / F_{500}}$ values than those observed for the total sample. If we consider galaxies with longer submm data (green and blue points) we also see that it is very difficult to distinguish, using these FIR/submm colours only, which galaxies will present an excess at longer wavelengths. Nonetheless galaxies combining high $F_{250} / F_{500}(\gtrsim 6.5)$ and high $F_{160} / F_{250}$ ratios ( $\gtrsim 2.0$ ), i.e., the most extreme cases, will show an excess at longer submm wavelengths.

If we now look at the top right panel of Fig. 8.14 where we consider the $F_{160} / F_{250}$ and the $F_{500} / F_{850(870)}$ ratios, we see that the $F_{500} / F_{850(870)}$ is the main driver of the excess at longer submm wavelengths. In our sample of 21 galaxies detected at 850 or $870 \mu \mathrm{~m}$, all of the galaxies having $F_{500} / F_{850(870)} \lesssim 5.5$ present an excess at wavelengths $\geq 850 \mu \mathrm{~m}$.


Fig. 8.14. (left) $F_{160} / F_{250}$ as a function of $F_{250} / F_{500}$ for the galaxies detected at $500 \mu \mathrm{~m}$. Stars and crosses differentiate between galaxies with or without excess at $500 \mu \mathrm{~m}$, respectively. The solid line denotes the median ratio of $\frac{F_{160} / F_{250}}{F_{250} / F_{500}} \sim 0.37$. (right) $F_{160} / F_{250}$ as a function of $F_{500} / F_{850(870)}$ for the galaxies detected at 850 or $870 \mu \mathrm{~m}$. Stars and crosses differentiate this time between galaxies with or without excess at 850(870) $\mu \mathrm{m}$, respectively. On the top panels, the colours indicate the status of the excess at longer wavelength: no data available (red), data available but no excess (blue), data available and excess (green) (same as in Fig. 8.12). On the bottom panels, the colours indicate the $\mathrm{S} / \mathrm{N}$ of the excess at $500 \mu \mathrm{~m}$ (bottom left) and 850(870) $\mu \mathrm{m}$ (bottom right).

The bottom panels of Fig. 8.14 shows the same colour-colour diagrams as in the top panels but this time colour coded by $\mathrm{S} / \mathrm{N}$ of the excess at 500 and $850(870) \mu \mathrm{m}$. We see that the $\mathrm{S} / \mathrm{N}$ of the excess is mainly driven by the submm colours $F_{250} / F_{500}$ for the $\mathrm{S} / \mathrm{N}$ at $500 \mu \mathrm{~m}$, and $F_{500} / F_{850(870)}$ for the $\mathrm{S} / \mathrm{N}$ at $850(870) \mu \mathrm{m}$. This is not surprising given that a lower $F_{250} / F_{500}$ or $F_{500} / F_{850(870)}$ ratio implies a shallower slope of the SED and thus a potentially stronger excess.

### 8.3.4 Explaining the excess: testing another dust composition

In this section we want to explore one of the hypotheses put forward to explain the submm excess in our galaxies: use another dust composition with grains more emissive in the submm. In our dust
model, we replace graphite ( Gr ) carbon grains by amorphous carbon (Ac) grains. As a consequence, the emissivity index of this dust composition is $\beta=1.7$, instead of 2.0 with our previous model: we show in Fig. 8.15 the comparison between the opacities of the two models.


Fig. 8.15. Comparison of the opacities, $\kappa_{a b s}$ for the two dust compositions: Gr (red) and Ac (blue). The $500 \mu \mathrm{~m}$ to $1000 \mu \mathrm{~m}$ zone is outlined in grey.

We regenerate the SEDs for our sub-sample of 78 galaxies using this alternative dust composition, and follow the same procedure to access to the residual at a given wavelength, $R(\lambda)_{A C}$, and the corresponding $\mathrm{S} / \mathrm{N}, \mathrm{S} / \mathrm{N}_{A C}=R(\lambda)_{A C} / \Delta R(\lambda)_{A C}$. The Ac model SEDs for the 37 galaxies with a submm excess from the Gr model, are shown in Fig. 8.20 along with the Gr model SEDs.

The new residual distributions at each wavelength are presented in Fig. 8.16. We see that at all wavelengths, the number of galaxies presenting an excess decreases but some galaxies still show an excess when modelled with amorphous carbon grains. The number of galaxies with a submm excess at $500 \mu \mathrm{~m}$ is now 14 , about half of the number of galaxies showing submm excess at $500 \mu \mathrm{~m}$ with the Gr model. At $850 / 870 \mu \mathrm{~m}$, the decrease in the presence of submm excess is less important: from 11 excess galaxies with graphite grains to eight with amorphous carbons. As for the four galaxies detected at $1200 \mu \mathrm{~m}$, two of them still present an excess with the Ac model. Thus in total, with the Ac model, 25 galaxies have a submm excess at least at one submm wavelength.

The $\mathrm{S} / \mathrm{N}$ of the excess in these galaxies is lower than those determined when using the Gr model (see Fig. 8.17): on average the $\mathrm{S} / \mathrm{N}$ has decreased by $27 \%, 31 \%$ and $53 \%$ at $500 \mu \mathrm{~m}, 850 / 870 \mu \mathrm{~m}$ and $1200 \mu \mathrm{~m}$ respectively. Fig. 8.17 also shows that galaxies with the flattest submm slopes, and the highest $\mathrm{S} / \mathrm{N}$ are the ones for which the excess remains with the Ac model: only part of the excess observed with the graphite grains model can be explained by the new dust composition.

Using more emissive dust grains - replacing graphite grains by amorphous carbons grains - can thus account for the submm excess observed in galaxies modelled with the Gr model for the weakest excesses. However, another mechanism needs to be invoked to explain the strongest submm excesses.

Fig. 8.18 shows where the submm excess galaxies lie for both models in the $\langle U\rangle$ vs $12+\log (\mathrm{O} / \mathrm{H})$ plane. We see that most galaxies for which the excess is explained by the amorphous carbon grains model are moderately metal-poor galaxies, with $8.2 \leq 12+\log (\mathrm{O} / \mathrm{H}) \leq 8.4$. This dust composition has already been shown to be more appropriate for LMC regions by Galliano et al. (2011) and Galametz et al. (2013), where the metallicity is $\sim 0.47 Z_{\odot}$ (as estimated by Galliano et al. 2011, i.e., $12+\log (\mathrm{O} / \mathrm{H})=8.36)$. Thus we believe that for galaxies with $12+\log (\mathrm{O} / \mathrm{H})$


Fig. 8.16. Distribution of the residuals at $500 \mu \mathrm{~m}$ (left), 850 and $870 \mu \mathrm{~m}$ (centre), and $1200 \mu \mathrm{~m}$ (right), for the DGS and KINGFISH galaxies. The KINGFISH distribution is shown on top of the DGS distribution. For all panels, the colours code the $\mathrm{S} / \mathrm{N}$ of the excess: $R(\lambda) / \Delta R(\lambda)$. A grey cell means a $\mathrm{S} / \mathrm{N}<1$ and thus no excess. The distribution of $R(\lambda)$ for the previous model using graphite grains is shown in light blue.


Fig. 8.17. $\mathrm{S} / \mathrm{N}_{G r}$ versus $\mathrm{S} / \mathrm{N}_{A C}$ for the three submm wavelengths 500 (top left), 850(870) (top right) and $1200 \mu \mathrm{~m}$ (bottom). For each panel, the symbols differentiate between galaxies without submm excess (crosses), with submm excess only with the graphite grains model ("Gr model", open diamonds) and with submm excess with both models ("Gr and Ac models", filled diamonds). The colour indicates the submm colour: $F_{250} / F_{500}$ (top left), $F_{500} / F_{850(870)}$ (top right), or $F_{500} / F_{1200}$ (bottom). Names are indicated for filled diamonds for the top right and bottom panel.


Fig. 8.18. Same as Fig.8.13 except that the symbols are now the same as in Fig. 8.17 (top left panel).
around 8.3 , the weakest submm excesses observed at $500 \mu \mathrm{~m}$ with a graphite grain model may be representative of the fact that a Galactic grain composition is not appropriate to model the dust emission of these galaxies.

## Impact on the dust mass



Fig. 8.19. Comparison of the dust mass obtained with the Ac model, $M_{d u s t} A C$, and the dust mass obtained with the Gr model, $M_{d u s t} G R$. The colours code the metallicity. The symbols are the same as in Fig. 8.17 top left panel. Names are indicated for galaxies where the Ac model can account for the submm excess.

Because the dust emissivity index is lower for the AC SED model, the grains are more emissive, thus requiring less dust mass to account for the same luminosity. And indeed the dust masses derived with the new dust composition are lower by a factor of $\sim 2.3\left(M_{\text {dust }} A C / M_{\text {dust }} G R \sim 0.43\right.$,
see Fig. 8.19). The Ac model SEDs for the 37 galaxies with a submm excess with the Gr model, are presented in Fig. 8.20 between $200 \mu \mathrm{~m}$ and $1300 \mu \mathrm{~m}$, together with the Gr model SEDs.

### 8.4 Summary

In this Chapter we saw that submm constraints combined to a realistic dust model is crucial to get an accurate estimate of the dust properties, and especially of the dust mass. Indeed a modified blackbody model tends to systematically underestimate the dust mass by a factor of 1.8 compared to a semi-empirical dust SED model. The temperature distribution in low-metallicity environments is broader and the dust is also globally warmer than in more metal-rich systems. The dust-to-stellar mass ratios is more or less constant over the metallicity range we are probing, but we found that, in this sample, dwarf galaxies are more efficient in forming stars than their metal-rich counterparts. We identified a number of galaxies presenting an excess at $500 \mu \mathrm{~m}$ and found that the submm excess appears preferentially in low-metallicity galaxies, but that the intensity of the excess does not depend on metallicity. An alternative dust composition with more emissive dust grains can fully explain the submm excess in about $1 / 3$ or the sample.


Fig. 8.20. SEDs modelled with the graphite grains model (red) and amorphous carbon grains (blue) for the galaxies showing a submm excess with the Gr model. To facilitate the comparison between the two modelled SEDs, we zoom into the $[200,1300] \mu \mathrm{m}$ range. The observed points are in black. The bottom panel of eqze plot indicates the residuals from the graphite grains model fit.


Fig. 8.20. Comparison of the Gr and Ac SEDs between $200 \mu \mathrm{~m}$ and $1300 \mu \mathrm{~m}$ (continued).


Fig. 8.20. Comparison of the Gr and Ac SEDs between $200 \mu \mathrm{~m}$ and $1300 \mu \mathrm{~m}$ (continued).


Fig. 8.20. Comparison of the Gr and Ac SEDs between $200 \mu \mathrm{~m}$ and $1300 \mu \mathrm{~m}$ (continued).


Fig. 8.20. Comparison of the Gr and Ac SEDs between $200 \mu \mathrm{~m}$ and $1300 \mu \mathrm{~m}$ (continued).

Table 8.1. Submm residuals for the DGS and KINGFISH samples in \%.

| Name | Exc BB ? | $\mathrm{R}(500) \pm \Delta \mathrm{R}(500)$ | Exc ? | $\mathrm{R}(850 / 870) \pm \Delta \mathrm{R}(850 / 870)$ | Exc ? | $\mathrm{R}(1200) \pm \Delta \mathrm{R}(1200)$ | Exc? |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| DGS |  |  |  |  |  |  |  |
| Haro11 | yes | $24.0 \pm 12.1$ | yes* | $235.6 \pm 32.4{ }^{L}$ | yes* |  |  |
| Haro2 |  | $-14.0 \pm 9.7$ |  | $9.3 \pm 37.4^{\text {S }}$ |  |  |  |
| Haro3 |  | $-12.2 \pm 9.4$ |  |  |  |  |  |
| He2-10 |  | $-13.3 \pm 8.6$ |  | $-16.8 \pm 12.9{ }^{S}$ |  | $-11.3 \pm 45.5$ |  |
| HS0052+2536 | yes | $113.8 \pm 50.8$ | yes* |  |  |  |  |
| IC10 |  | $-8.2 \pm 31.1$ |  |  |  |  |  |
| IIZw40 |  | $-2.2 \pm 8.5$ |  | $206.6 \pm 18.1{ }^{\text {S }}$ | yes* | $263.0 \pm 47.8$ | yes* |
| Mrk1089 |  | $-11.1 \pm 10.8$ |  | $76.7 \pm 14.5{ }^{L}$ | yes* |  |  |
| Mrk930 |  | $4.5 \pm 13.6$ |  | $428.5 \pm 14.7{ }^{L}$ | yes* |  |  |
| NGC1140 |  | $-11.3 \pm 10.8$ |  | $27.0 \pm 45.8{ }^{\text {S }}$ |  |  |  |
| NGC1569 |  | $-2.3 \pm 8.6$ |  | $35.9 \pm 20.8{ }^{S}$ | yes | $202.4 \pm 69.0$ | yes* |
| NGC1705 | yes | $-4.7 \pm 13.8$ |  | $589.9 \pm 68.7{ }^{L}$ | yes* |  |  |
| NGC2366 | yes | $12.8 \pm 9.7$ | yes |  |  |  |  |
| NGC4214 | yes | $0.9 \pm 8.2$ |  |  |  | $64.4 \pm 32.5$ | yes |
| NGC4449 |  | $-8.7 \pm 8.1$ |  |  |  |  |  |
| NGC4861 | yes | $-4.0 \pm 12.4$ |  |  |  |  |  |
| NGC5253 |  | $1.7 \pm 8.2$ |  | $171.2 \pm 17.1^{L}$ | yes* |  |  |
| NGC625 | yes | $3.3 \pm 8.6$ |  |  |  |  |  |
| NGC6822 | yes | $9.9 \pm 31.1$ |  |  |  |  |  |
| UM311 |  | $-10.5 \pm 9.0$ |  |  |  |  |  |
| UM448 |  | $3.8 \pm 9.8$ |  | $232.7 \pm 16.4{ }^{L}$ | yes* |  |  |
| VIIZw403 | yes | $60.8 \pm 32.7$ | yes* |  |  |  |  |
| KINGFISH |  |  |  |  |  |  |  |
| NGC0337 | yes | $16.4 \pm 8.3$ | yes | $28.1 \pm 18.0{ }^{S}$ | yes |  |  |
| NGC0628 |  | $6.3 \pm 8.4$ |  |  |  |  |  |
| NGC0855 |  | $-16.6 \pm 16.8$ |  |  |  |  |  |
| NGC0925 | yes | $30.5 \pm 8.0$ | yes* |  |  |  |  |
| NGC1097 |  | $1.4 \pm 8.0$ |  | $-37.5 \pm 71.5{ }^{\text {S }}$ |  |  |  |
| NGC1266 |  | $-4.4 \pm 11.5$ |  |  |  |  |  |
| NGC1291 |  | $5.4 \pm 9.2$ |  |  |  |  |  |
| NGC1316 |  | $6.3 \pm 31.3$ |  |  |  |  |  |
| NGC1377 | yes | $24.4 \pm 17.9$ | yes |  |  |  |  |
| IC0342 |  | $-3.4 \pm 7.6$ |  |  |  |  |  |
| NGC1482 |  | $-5.6 \pm 8.6$ |  | $-36.9 \pm 27.5^{S}$ |  |  |  |
| NGC1512 | yes | $12.8 \pm 9.0$ | yes |  |  |  |  |
| NGC2146 |  | $-10.5 \pm 7.9$ |  |  |  |  |  |
| HoII |  | $8.1 \pm 55.9$ |  |  |  |  |  |
| NGC2798 |  | $3.1 \pm 9.3$ |  | $-6.7 \pm 25.1{ }^{S}$ |  |  |  |
| NGC2841 |  | $1.3 \pm 7.9$ |  |  |  |  |  |
| NGC2915 | yes | $14.2 \pm 14.0$ | yes |  |  |  |  |
| HoI | yes | $25.4 \pm 33.8$ |  |  |  |  |  |
| NGC2976 |  | $7.3 \pm 8.1$ |  | $-11.0 \pm 43.1{ }^{\text {S }}$ |  |  |  |
| NGC3049 | yes | $48.7 \pm 9.1$ | yes* |  |  |  |  |
| NGC3077 | yes | $20.5 \pm 8.5$ | yes* |  |  |  |  |
| M81dwB |  | $33.7 \pm 55.7$ |  |  |  |  |  |
| NGC3190 |  | $-8.1 \pm 8.9$ |  | $-23.7 \pm 24.8{ }^{\text {S }}$ |  |  |  |
| NGC3184 |  | $8.3 \pm 8.1$ | yes |  |  |  |  |
| NGC3198 | yes | $17.0 \pm 8.0$ | yes |  |  |  |  |
| IC2574 | yes | $-0.3 \pm 9.9$ |  |  |  |  |  |
| NGC3265 | yes | $23.2 \pm 16.4$ | yes |  |  |  |  |
| NGC3351 |  | $-6.9 \pm 8.2$ |  |  |  |  |  |
| NGC3521 |  | $4.8 \pm 8.0$ |  | $-31.9 \pm 47.4{ }^{S}$ |  |  |  |
| NGC3621 | yes | $17.1 \pm 7.6$ | yes |  |  |  |  |
| NGC3627 |  | $-6.6 \pm 7.7$ |  | $-38.7 \pm 54.2{ }^{\text {S }}$ |  |  |  |
| NGC3773 |  | $6.0 \pm 14.9$ |  |  |  |  |  |
| NGC3938 |  | $5.2 \pm 8.2$ |  |  |  |  |  |
| NGC4236 | yes | $44.8 \pm 8.5$ | yes* |  |  |  |  |
| NGC4254 |  | $-10.5 \pm 8.0$ |  | $-56.0 \pm 89.4^{S}$ |  |  |  |
| NGC4321 |  | $-2.7 \pm 8.3$ |  | $-63.3 \pm 87.0{ }^{\text {S }}$ |  |  |  |
| NGC4536 | yes | $15.0 \pm 8.1$ | yes | $-56.7 \pm 38.6^{S}$ |  |  |  |
| NGC4559 | yes | $26.6 \pm 8.0$ | yes* |  |  |  |  |
| NGC4569 |  | $-10.0 \pm 8.2$ |  | $-35.9 \pm 21.2{ }^{\text {S }}$ |  |  |  |
| NGC4579 |  | $-3.2 \pm 8.2$ |  | $-29.1 \pm 19.2{ }^{\text {S }}$ |  |  |  |
| NGC4594 | yes | $-0.6 \pm 9.5$ |  | $-66.8 \pm 80.3^{S}$ |  |  |  |
| NGC4625 | yes | $16.9 \pm 10.7$ | yes |  |  |  |  |

Table 8.1. (continued) Submm residuals for the DGS and KINGFISH samples in \%.

| Name | Exc BB ? | $\mathrm{R}(500) \pm \Delta \mathrm{R}(500)$ | Exc ? | $\mathrm{R}(850 / 870) \pm \Delta \mathrm{R}(850 / 870)$ | Exc ? | $\mathrm{R}(1200) \pm \Delta \mathrm{R}(1200)$ | Exc ? |
| :--- | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| NGC4631 | yes | $14.6 \pm 7.9$ | yes* | $70.5 \pm 24.8^{\mathrm{S}}$ | yes* |  |  |
| NGC4725 |  | $12.6 \pm 8.0$ | yes |  |  |  |  |
| NGC4736 | yes | $14.6 \pm 8.5$ | yes | $-6.1 \pm 52.1 \mathrm{~S}$ |  |  |  |
| NGC4826 |  | $-2.5 \pm 8.1$ |  | $11.5 \pm 40.8^{S}$ |  |  |  |
| NGC5055 |  | $0.8 \pm 8.1$ |  |  |  |  |  |
| NGC5398 | yes | $19.1 \pm 9.9$ | yes* |  |  |  |  |
| NGC5408 | yes | $61.2 \pm 22.4$ | yes* |  | yes |  |  |
| NGC5457 | yes | $18.2 \pm 7.9$ | yes* |  |  |  |  |
| NGC5474 | yes | $22.1 \pm 9.2$ | yes* |  |  |  |  |
| NGC5713 |  | $-8.1 \pm 8.2$ |  | $42.7 \pm 25.6^{S}$ | $-28.3 \pm 29.3 \mathrm{~S}$ |  |  |
| NGC5866 |  | $-5.5 \pm 9.8$ |  | $-74.9 \pm 28.5 \mathrm{~S}$ |  |  |  |
| NGC6946 |  | $-6.7 \pm 7.8$ |  | $-18.6 \pm 21.4 S$ |  |  |  |
| NGC7331 |  | $4.2 \pm 8.1$ |  |  |  |  |  |
| NGC7793 | yes | $24.9 \pm 8.0$ | yes* |  |  |  |  |

${ }^{L}$ LABOCA residual at $870 \mu \mathrm{~m}$.
${ }^{S}$ SCUBA residual at $850 \mu \mathrm{~m}$.

* Excess remains if use amorphous carbon grains model.


## Chapter 9

## Gas-to-Dust mass ratios as a function of metallicity over a 2 dex metallicity range

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In this last Chapter, we incorporate the gas into our dust picture, and look at the gas-to-dust mass ratios (G/D) as a function of metallicity. The G/D links the amount of metals locked up in dust and in the gas phase and is thus a powerful tracer of the evolutionary stage of a galaxy. Investigations of the relation between the observed G/D and metallicity can thus place important constraints on the physical processes governing galaxy evolution and more specifically on chemical evolution models.

### 9.1 Motivations

The observed G/D of integrated galaxies as a function of metallicity has been intensively studied over the past decades (e.g., Issa et al. 1990; Lisenfeld \& Ferrara 1998; Hirashita et al. 2002; James et al. 2002; Hunt et al. 2005; Draine et al. 2007; Engelbracht et al. 2008; Galliano et al. 2008; Muñoz-Mateos et al. 2009; Bendo et al. 2010b; Galametz et al. 2011; Magrini et al. 2011). In the disk of our Galaxy the proportion of heavy elements in the gas and in the dust has been shown to scale with the metallicity (Dwek 1998) if one assumes that the time dependence of the dust formation timescale is the same as that of the dust destruction timescale. This results in a constant dust-to-metal mass ratio and gives a dependence of the G/D on metallicity as: G/D $\propto \mathrm{Z}^{-1}$ (that we call hereafter the "reference" trend). This reference trend between G/D and metallicity seems consistent with the observations of galaxies with near-solar metallicities (e.g., James et al. 2002; Draine et al. 2007; Bendo et al. 2010b; Magrini et al. 2011). However some studies also show that the G/D of some low-metallicity dwarf galaxies deviate from this reference trend, with a higher G/D than expected for their metallicity (Lisenfeld \& Ferrara 1998; Galliano et al. 2003; Hunt et al. 2005; Galliano et al. 2005, 2008; Bernard et al. 2008; Engelbracht et al. 2008; Galametz et al. 2011; Galliano et al. 2011).

The G/D is often used to empirically estimate the "CO-free" molecular gas. This molecular gas, not directly traced by CO measurements, was first proposed in low-metallicity galaxies using the $158 \mu \mathrm{~m}$ [CiI] line (Poglitsch et al. 1995; Israel et al. 1996; Madden et al. 1997, 2012). Alternatively, using the dust mass determined from FIR measurements and assuming a G/D, given the metallicity of the galaxy, a gas mass can be estimated. This gas mass is then compared to the observed Hi and CO gas masses to estimate the "CO-free" gas mass. This method is also used to estimate the CO-to-H2 conversion factor in local (e.g., Guelin et al. 1993, 1995; Neininger et al. 1996; Boselli et al. 2002; Sandstrom et al. 2013) and high- $z$ galaxies (e.g., Magdis et al. 2012; Magnelli et al. 2012). However using these methods requires an accurate estimation of the G/D at a given metallicity.

A certain number of instrumental limitations and/or model caveats have limited former studies of the G/D. First, limits in wavelength coverage in the FIR have hampered the precise determination of the dust masses. For the earliest studies, the dust masses were derived from IRAS or Spitzer measurements, not extending further than $100-160 \mu \mathrm{~m}$, therefore not tracing the cooler dust. As the bulk of the dust mass in galaxies often resides in the cold dust component, this has important consequences for the dust mass and therefore G/D determination. Before Herschel, using Spitzer and ground-based data, Galametz et al. (2011) indeed showed that a broad wavelength coverage of the FIR-to-submm part of the SED was critical to obtain accurate estimates of the dust masses. We confirmed these results in Chapter 8. Second, some of the studies presented previously used modified blackbody models to derive dust masses. Using Herschel data and a semi-empirical dust model, we showed in Chapter 8 that the dust mass modelled with a modified blackbody could be underestimated by a factor of $\sim 2$, consistent with the findings of Dale et al. (2012) on the KINGFISH sample. And finally, the limited sensitivities of the pre-Herschel era instruments only allowed the detection of dust for the brightest and highest-metallicity dwarf galaxies, limiting G/D studies to metallicities $\geq 1 / 5 Z_{\odot}(12+\log (\mathrm{O} / \mathrm{H})=8.0)$.

In this work we investigate the relation between the G/D and metallicity avoiding the limitations
and caveats mentioned in the previous paragraph. Using our new Herschel data to constrain a semi-empirical dust model, we have a more accurate determination of the dust masses compared to Spitzer-only dust masses and/or modified blackbody dust masses. Our sample also covers a wide range in metallicity ( 2 dex, from $12+\log (\mathrm{O} / \mathrm{H})=7.1$ to 9.1 ), with a significant fraction of the sample below $12+\log (\mathrm{O} / \mathrm{H})=8.0(\sim 30 \%$, see Section 9.2), thanks to the increased sensitivity of Herschel which enables us to access the dust in the lowest metallicity galaxies. We are thus able to provide better constraints on the G/D at low-metallicities.

In Section 9.2, we describe the sample and the method used to estimate dust and total gas masses are described in Section 9.3 and Section 9.4 respectively. Then we investigate the relation of the observed G/D with metallicity in Section 9.5 and fit several empirical relations to the data. We then interpret our results with the aid of several chemical evolution models in Section 9.6. Throughout we consider (G/D) $\odot=162^{1}$ (Zubko et al. 2004).

### 9.2 Sample

We combine 3 different samples for our study of the G/D: the Dwarf Galaxy Survey (DGS, Madden et al. 2013), the KINGFISH survey (Kennicutt et al. 2011) and a subsample of the sample presented in Galametz et al. (2011) (called the "G11 sample" hereafter). The basic parameters for all of the galaxies such as positions and distances can be found in Madden et al. (2013) and Chapter 3 for the DGS, Kennicutt et al. (2011) for KINGFISH and Galametz et al. (2011) for the G11 sample.

The G11 sample consists of all of the galaxies in Galametz et al. (2011) that are neither already in the DGS nor in the KINGFISH samples, except those galaxies which show a submm excess (see Section 9.3). This gives 17 additional galaxies, mostly solar or super-solar environments (mostly spiral galaxies), with metallicities from $12+\log (\mathrm{O} / \mathrm{H})=8.14$ to 9.1 . The metallicity distribution for each of the 3 samples is presented in Fig. 9.1.

All of these metallicities have been derived using empirical strong emission line methods (see Madden et al. 2013 for the DGS, Kennicutt et al. 2011 for KINGFISH and Galametz et al. 2011 for G11 metallicity determination). The DGS and KINGFISH metallicities have been obtained through the $\mathrm{R}_{23}$ ratio $^{2}$ with the Pilyugin \& Thuan (2005) calibration (see Chapter 3). Galametz et al. (2011) do not indicate precisely which calibration they use to convert $\mathrm{R}_{23}$ into metallicity, thus several metallicities for the G11 galaxies were re-estimated from the original line intensities with the Pilyugin \& Thuan (2005) calibration. We also assume a conservative 0.1 dex uncertainty for the G11 metallicities. On average for the total sample, the uncertainty on the metallicity measurements is $\sim 0.1$ dex. The metallicities for the whole sample are listed in Table 9.3 . Other methods exist to determine metallicities and can lead to very different values, but this will only introduce a systematic offset in the adopted values here (Kewley \& Ellison 2008). Note that these metallicity values correspond to global estimates. On smaller scales within galaxies, differences can occur due to inhomogeneous mixing of metals: metallicity gradients have been observed in large spiral galaxies (Garnett et al. 2004; Bendo et al. 2010b; Moustakas et al. 2010). Dwarf galaxies, however, are smaller in size than metal-rich galaxies and we can presume, for this study, that metallicity is more homogeneous within these environments (Revaz \& Jablonka 2012; Valcke et al. 2008) .

This gives us a total of 126 galaxies spanning a 2 dex range in metallicity (Fig. 9.1). We see that the low-metallicity end of the distribution is fairly well sampled as we have $\sim 30 \%$ of the total sample with metallicities below $1 / 5 Z_{\odot}$.

[^47]

Fig. 9.1. Metallicity distribution of the DGS (purple), KINGFISH (orange) and G11 (green) samples from $12+\log (\mathrm{O} / \mathrm{H})=7.14$ to 9.1 . The total distribution is indicated in grey. Solar metallicity is indicated here as a guide to the eye, as well as the $1 / 50,1 / 10,1 / 5$ and $2 Z_{\odot}$ values.

### 9.3 Dust masses

To ensure a consistent determination of the dust masses throughout our sample, all of the galaxies are modelled with the dust SED model presented in Galliano et al. (2011). The dust masses for the DGS and KINGFISH samples can be found in Chapter 7. A submm excess is observed in some DGS and KINGFISH galaxies at $500 \mu \mathrm{~m}$ (Chapter 8, Dale et al. 2012). If present, the excess is rather small at $500 \mu \mathrm{~m}$ (the median $\mathrm{S} / \mathrm{N}$ of the excess at $500 \mu \mathrm{~m}$ is $\sim 2.0$ ) and can increase as we go to longer wavelengths (see Chapter 8). However, because of the unknown origin of this excess and because of the uncertainties it can bring in the dust mass estimation, we do not attempt to model this excess with additional modifications to the model. We discuss this choice in Section 9.5.3.

We also model galaxies from the G11 sample, as some model assumptions were different in Galametz et al. (2011). Herschel constraints are not present but other submm constraints are taken into account such as JCMT/SCUBA at $850 \mu \mathrm{~m}$ and/or APEX/LABOCA at $870 \mu \mathrm{~m}$, allowing a precise determination of the dust masses (given in Table 9.1). Galametz et al. (2011) also observed a submm excess in 9 galaxies of their original sample and modelled it with a Very Cold Dust (VCD) component. However the submm excess is not fully understood yet and this extra VCD component may lead to an overestimation of the dust mass. Because we do not have constraints between 160 $\mu \mathrm{m}$ and the available ground-based submm fluxes to see where the submm excess starts, we do not consider nor model these galaxies here.

The wavelength coverage is not exactly the same from galaxy to galaxy. The most important constraints for the determination of the dust mass are constraints sampling the peak of the dust SED. Herschel provides such constraints for all of the DGS and KINGFISH galaxies. Some dwarf galaxies are not detected with Herschel at submm wavelengths (beyond $250 \mu \mathrm{~m}$ ) and are noted in Fig. 9.4 (see Section 9.4). These galaxies harbour particularly warm dust (see Chapter 8, RémyRuyer et al. 2013); the peak of their SED is thus shifted towards shorter wavelengths and is then well sampled by constraints until $160 \mu \mathrm{~m}$ where the galaxy is still detected. For galaxies in the G11 sample, the peak of the dust SED is probed by Spitzer observations and the Rayleigh Jeans slope of the SED by longer submm wavelength observations. Thus we are confident in the dust masses
we derive with these sets of constraints.
The errors on the dust masses for the G11 sample are estimated following the same procedure as for the DGS and KINGFISH galaxies in Chapter 7, by generating 300 random realisations of the SED, perturbed according to the random and systematic noise, in order to get a distribution for the dust mass. The error bars on the dust mass are taken to be the $66.67 \%$ confidence interval of the distribution (i.e., the range of the parameter values between 0.1667 and 0.8333 of the repartition function). The dust masses and uncertainties for the G11 sample are presented in Table 9.1.

Table 9.1. Dust masses for the G11 sample.

| Name | $\mathrm{M}_{\text {dust }}\left[M_{\odot}\right]$ | Uncertainty (\%) |
| :--- | :---: | :---: |
| M83 | $1.10 \times 10^{7}$ | 8.14 |
| NGC1808 | $2.54 \times 10^{7}$ | 34.24 |
| NGC7552 | $7.24 \times 10^{7}$ | 16.10 |
| M82 | $6.85 \times 10^{6}$ | 33.02 |
| NGC1068 | $2.77 \times 10^{8}$ | 36.05 |
| NGC0891 | $5.84 \times 10^{7}$ | 11.84 |
| MGC+02-04-025 | $1.01 \times 10^{8}$ | 47.08 |
| NGC7469 | $1.55 \times 10^{8}$ | 19.16 |
| NGC5256 | $2.01 \times 10^{8}$ | 62.93 |
| NGC5953 | $2.25 \times 10^{7}$ | 18.90 |
| M51 | $3.74 \times 10^{8}$ | 21.82 |
| NGC3995 | $5.75 \times 10^{7}$ | 23.47 |
| NGC3994 | $5.27 \times 10^{7}$ | 34.81 |
| NGC6052 | $5.84 \times 10^{7}$ | 32.58 |
| NGC1222 | $1.33 \times 10^{7}$ | 26.71 |
| NGC7674 | $9.83 \times 10^{7}$ | 29.69 |
| NGC4670 | $1.13 \times 10^{6}$ | 21.12 |

### 9.4 Gas masses from the literature

### 9.4.1 Atomic gas masses

The Hi masses and their errors are compiled from the literature, and rescaled to the distances used here. Most of the atomic gas masses are given in Galametz et al. (2011) for the G11 sample, and in Draine \& Li (2007) for the KINGFISH survey. They are presented in Madden et al. (2013) and Chapter 3 for the DGS. The errors were not available for all of the Hi measurements. When no error was available for the Hi mass, we adopted the mean value of all of the relative errors on the Hi masses compiled from the literature: $\sim 16 \%$.

## Correcting the HI masses

However the Hi extent of a galaxy is not necessarily the same as the aperture used to probe the dust SED, as the Hi often extends beyond the optical radius of a galaxy (Hunter 1997). This can be particularly true for dwarf galaxies where the Hi halo can be very extended: some irregular galaxies present unusually extended Hi gas (up to seven times the optical radius, Huchtmeier 1979; Huchtmeier et al. 1981; Carignan \& Beaulieu 1989; Carignan et al. 1990; Thuan et al. 2004). We also note that in some galaxies (e.g., NGC 4449), gas morphology may be highly perturbed due to past interactions or mergers (e.g., Hunter et al. 1999). This may also lead to significant uncertainties in the Hi mass and thus on the derived G/D (e.g., Karczewski et al. 2013). Thus we check the literature for the DGS sample for the size of the Hı halo to compare it to the dust aperture. It was not possible to find this information for $\sim 38 \%$ of the sample (Hi not detected or no map available).


Fig. 9.2. G/D mass ratios as a function of metallicity for the 2 values of $X_{C O} \quad X_{C O, M W}$ (right) and $X_{C O, Z}$ (left), before (top) and after (bottom) correcting the HI masses. The mean error for the datapoints is shown in grey. On both panels, the colours delineate the three samples: DGS in purple, KINGFISH in orange and G11 in green. On the bottom panels, the symbol traces the changes made in the HI masses (see text for details on the uncertainties and the changes on the atomic gas masses).

For the rest, $25 \%$ of the DGS galaxies have a Hi extent that corresponds to the dust IR aperture, which has been chosen to be 1.5 times the optical radius (for most cases, see Chapter 4 and RémyRuyer et al. 2013); and $35 \%$ have a Hi halo that is more extended. If we assume that the HI mass distribution follows an exponential profile (based on the observed high central gas concentration seen in BCDs, e.g., van Zee et al. 1998, 2001; Simpson \& Gottesman 2000), we can correct the total Hi mass for these galaxies to find the Hı mass corresponding to the dust aperture.

$$
\begin{equation*}
M_{H I}(\text { IRcorr })=M_{H I} \times \frac{\int_{0}^{r_{\text {dust }}} \exp ^{-r^{2} /\left(2 \sigma^{2}\right)} \mathrm{d} r}{\int_{0}^{r_{H I}} \exp ^{-r^{2} /\left(2 \sigma^{2}\right)} \mathrm{d} r} \tag{9.1}
\end{equation*}
$$

where $r_{\text {dust }}$ and $r_{H I}$ are the dust and Hi radii respectively, and $\sigma$ the standard deviation of the gaussian profile: $\sigma=r_{H I} / 2.35$.

In reality, the Hi profile can show a complicated structure with clumps and shells, rendering the profile more assymetric. Our correction corresponds to a factor of $\sim 1.55$ on average, for these galaxies. Several studies (Thuan \& Martin 1981; Swaters et al. 2002; Lee et al. 2002; Begum \& Chengalur 2005a; Pustilnik \& Martin 2007) have tried to quantify the extent of the Hi halo for dwarf galaxies and found that the ratio of Hi size to the optical size is typically 2 , which gives a
correction of $\sim 1.4$. This is similar to what we find here with our assumptions. The atomic gas masses for the sample, after correction if needed, are presented in Table 9.3. The G/D mass ratio before and after this correction is shown in Fig. 9.2.

### 9.4.2 Molecular gas masses

The $\mathrm{H}_{2}$ masses have been compiled from the literature. They have been rescaled, when necessary, to the distance adopted here to derive the dust masses.

There are two main issues in determining the molecular gas masses. First, detection of CO in low-metallicity galaxies is challenging: sensitivity has limited CO detections to galaxies with $12+\log (\mathrm{O} / \mathrm{H}) \gtrsim 8.0$ (e.g., Leroy et al. 2009; Schruba et al. 2012). The other issue in the $\mathrm{H}_{2}$ mass determination is the choice of the conversion factor between CO intensities and molecular gas masses, $X_{C O}$. Indeed the variation of this factor with metallicity is poorly constrained, and a number of studies have been dedicated to quantifying the dependence of $X_{C O}$ on metallicity (Wilson 1995; Israel 1997; Boselli et al. 2002; Israel et al. 2003; Strong et al. 2004; Leroy et al. 2011; Schruba et al. 2012; Bolatto et al. 2013). From a sample of 16 dwarf galaxies, and assuming a constant $\mathrm{H}_{2}$ depletion timescale, Schruba et al. (2012) found a $X_{C O}$ scaling with $(\mathrm{O} / \mathrm{H})^{-2}$. This relation takes into account possible "CO-free" gas as the $X_{C O}$ conversion factor is estimated from the total reservoir of molecular gas needed for star formation (Schruba et al. 2012). Following Cormier et al. (2014), we estimate the molecular gas masses from a constant $X_{C O}$ factor using the Galactic value: $X_{C O, M W}=2.0 \times 10^{20} \mathrm{~cm}^{-2}\left(\mathrm{~K} \mathrm{~km} \mathrm{~s}^{-1}\right)^{-1}$ (Ackermann et al. 2011), giving us $M_{H_{2}, M W}$, and from a $X_{C O}$ factor depending on $(\mathrm{O} / \mathrm{H})^{-2}: X_{C O, Z}$, giving us $M_{H_{2}, Z}$ :

$$
\begin{equation*}
X_{C O, Z}=X_{C O, M W} \times\left(\frac{Z_{\odot}}{Z}\right)^{2} \tag{9.2}
\end{equation*}
$$

This provides a conservative range of molecular gas mass estimates that reflects how uncertain the molecular gas mass determination is. For this reason, we do not give any error bar on our molecular gas mass.

## Correcting the $\mathbf{H}_{2}$ masses

In order to go beyond the CO upper limits and to constrain the G/D behaviour at low metallicities we find a way to estimate the amount of molecular gas for the lowest metallicity galaxies. Figure 9.3 shows the ratio of $M_{H_{2}}$-to- $M_{H I}$ as a function of metallicity for our sample and for both cases of $X_{C O}$. We note that around $12+\log (\mathrm{O} / \mathrm{H}) \sim 8.1$ the ratio $M_{H_{2}} / M_{H I}$ drops suddenly for the detected galaxies for both $X_{C O}$. For these very low-metallicity galaxies with $12+\log (\mathrm{O} / \mathrm{H}) \leq 8.1$, the mean ratio between the detected $M_{H_{2}}$ and $M_{H I}$ is $1.2 \%$, for $X_{C O, M W}$. Using $X_{C O, Z}$, this ratio goes up to $68 \%$, the molecular gas mass is not negligible anymore but is still more than a factor of 2 lower than the atomic gas mass. Thus for galaxies with non-detections in CO, or without any CO observations, and with $12+\log (\mathrm{O} / \mathrm{H}) \leq 8.1$, we replace the upper limit values by $0.012 \times M_{H I}$ for $M_{H_{2}, M W}$ and $0.68 \times M_{H I}$ for $M_{H_{2}, Z}$. Given the low molecular gas fraction we find, this will not greatly affect our interpretation of G/D nor the conclusions. From now on, the galaxies for which we apply this correction will be treated as detections. This $12+\log (\mathrm{O} / \mathrm{H})$ value of $\sim 8.1$ has already been noted as being special for dwarf galaxies (e.g., for the strength of the PAH features: Engelbracht et al. 2005; Madden et al. 2006; Draine et al. 2007; Engelbracht et al. 2008; Galliano et al. 2008). The molecular gas masses we use in the following analysis are presented in Table 9.3. The difference between the G/D before and after the $\mathrm{H}_{2}$ mass correction can be seen by comparing Fig. 9.2 bottom and Fig. 9.4.


Fig. 9.3. $M_{H_{2}} / M_{H I}$ as a function of metallicity for the whole sample. The blue crosses are for molecular gas masses computed with $X_{C O, M W}$ and the orange diamonds are for molecular gas masses computed with $X_{C O, Z}$. Upper limits in the molecular gas mass are indicated with grey arrows and smaller grey symbols. The mean error for the data points is shown in grey on the bottom right of the plot. The plain line shows the unity line. The dashed blue and orange lines show the $1.2 \%$ and $68 \%$ molecular-to-atomic gas mass fractions respectively and represent the mean $\mathrm{H}_{2}$-to- HI ratio of the detected galaxies with $12+\log (\mathrm{O} / \mathrm{H})<8.1$ (see text).

### 9.4.3 Total gas masses

We get the total gas mass, $M_{g a s}$, by adding all of the different gas contributions: the atomic gas mass, the molecular gas mass, the helium gas mass and the gaseous metal mass:

$$
\begin{equation*}
M_{g a s}=M_{H I}(\text { IRcorr })+M_{H_{2}}+M_{H e}+Z \times M_{g a s}, \tag{9.3}
\end{equation*}
$$

where $M_{\mathrm{He}}$ is the helium mass and Z the mass fraction of metals in the galaxy. Assuming $M_{\mathrm{He}}=Y_{\odot} M_{\mathrm{gas}}$, where $Y_{\odot}$ is the Galactic mass fraction of Helium, $Y_{\odot}=0.270$ (Asplund et al. 2009), we have:

$$
\begin{equation*}
M_{g a s}=\mu_{g a l}\left(M_{H I}+M_{H_{2}}\right), \tag{9.4}
\end{equation*}
$$

with $\mu_{g a l}=1 /\left(1-Y_{\odot}-\mathrm{Z}\right)$ the mean atomic weight. $\mu_{g a l}$ has been computed for each galaxy and the mean value for our sample is $1.38 \pm 0.01$ (see Table 9.3). We get $Z$ assuming $\left(\mathrm{Z} / Z_{\odot}\right)=(\mathrm{O} / \mathrm{H}) /\left(\mathrm{O} / \mathrm{H}_{\odot}\right)$ and $Z_{\odot}=0.014$ (Asplund et al. 2009).

We assume here that the ionised gas mass ( $M_{\mathrm{HII}}$ ) is negligible compared to the Hi mass. We perform the test for 67 galaxies of the sample, with $M_{\text {HII }}$ derived from $\mathrm{H} \alpha$ measurements of Gil de Paz et al. (2003); Kennicutt et al. (2009); Skibba et al. (2011) and found $M_{\mathrm{HII}} / M_{H I} \sim 0.2 \%$. However, we found two dwarf galaxies for which the ionised gas mass should be taken into account as it contributes equally or more than the atomic gas mass: Haro11 ( $M_{\mathrm{HII}} \sim 1.2 \times M_{H I}$, Cormier et al. 2012) and Pox186 ( $M_{\mathrm{HII}} \sim M_{H I}$, Gil de Paz et al. 2003). For these two galaxies, the total gas mass also includes $M_{\text {HII }}$. For Haro11 the ionised gas mass is taken from Cormier et al. (2012). For Pox186, we assume that the mass of ionised gas $M_{H I I}$ is the mass coming from the $\mathrm{H} \alpha$ emission and is estimated from the $\mathrm{H} \alpha$ luminosity and the formula of Fathi et al. (2011):

$$
\begin{equation*}
M_{H \alpha}=m_{H} \times \frac{L_{H \alpha}}{3.56 \times 10^{-25} n_{e}} \tag{9.5}
\end{equation*}
$$

where $m_{H}$ is the mass of the hydrogen proton in $M_{\odot}, L_{H \alpha}$ is the total $\mathrm{H} \alpha$ luminosity in erg. $\mathrm{s}^{-1}$, and $n_{e}$ is the electron number density in $\mathrm{cm}^{-3} . n_{e}=70 \mathrm{~cm}^{-3}$ in the HII region for Pox186 (Guseva et al. 2004). For these two galaxies, the total gas mass also includes $M_{\text {HII }}$.

The G/D as a function of metallicity is presented in Fig. 9.4 a bb for the two cases: $X_{C O, M W}$ or $X_{C O, Z}$. The average error on the observed G/D is $\sim 27 \%$ in both $X_{C O}$ cases ( $\sim 10 \%$ for the total gas mass and $\sim 26 \%$ for the dust mass). The dashed line indicates the reference scaling of the G/D with metallicity. The colours of the symbols indicate the reliability of the data points by tracing if the gas or dust masses determinations are uncertain. Blue symbols refer to Hz or $\mathrm{H}_{2}$ non-detections or to the absence of $\mathrm{H}_{2}$ observations for the galaxy. Red symbols indicate that the galaxy is not detected at wavelengths $\geq 160 \mu \mathrm{~m}$. The combination of both indications for the gas and dust masses is shown with the purple symbols. Black symbols indicate that both gas and dust masses have reliable measurements ( $67 \%$ of the sample).

The type of symbols indicates whether or not the HI and/or $\mathrm{H}_{2}$ masses have been corrected. For the Hi masses we distinguish three cases for the DGS galaxies: the Hi extent of the galaxy is unknown and we cannot correct the Hi mass (diamonds), the Hi extent is known and greater than the dust aperture and we correct the Hi mass (triangles) and the Hi extent is known and similar to the dust aperture, there is no need to correct the Hi mass (crosses). The galaxies with $12+\log (\mathrm{O} / \mathrm{H})<8.1$ for which the $\mathrm{H}_{2}$ masses have been corrected (either from upper limit or lack of measurements) are indicated as filled symbols (see Section 9.4.2 on $\mathrm{H}_{2}$ masses).

### 9.5 Analysis

### 9.5.1 Observed gas-to-dust mass ratio - metallicity relation and dispersion

To evaluate the general behaviour and scatter in the G/D values at different metallicities, we consider the error-weighted mean values of $\log (\mathrm{G} / \mathrm{D})$ in metallicity bins (neglecting the upper/lower limits), with the bin sizes chosen to include at least two galaxies and to span at least 0.1 dex. The result is overlaid as pink filled circles in Fig. 9.4 c\&d. We also look at the dispersion of the G/D values in each metallicity bin (see bottom panels of Fig. $9.4 \mathrm{c} \mathrm{\& d}$ ), by computing the standard deviation of the $\log (\mathrm{G} / \mathrm{D})$ values in each bin (also neglecting the upper/lower limits). The dispersion is $\sim$ 0.37 dex (i.e., a factor of 2.3 ) on average for all bins and for both $X_{C O}$ values. Additionally, in one bin the G/D vary on average by one order of magnitude. This confirms that the relation between G/D and metallicity is not trivial even at a given metallicity, over the whole metallicity range. We also see that the dispersion in the observed G/D values does not depend on metallicity. This indicates that the scatter within each bin may be intrinsic and does not reflect systematic observational or correction errors. This also means that the metallicity is not the only driver for the observed scatter in the G/D values: other processes operating in galaxies can lead to large variation in the G/D in a given metallicity range, throughout this range. However, there might be a selection bias in our sample. Indeed our sample is mainly composed of star-forming gas-rich dwarf galaxies at low metallicities and spiral galaxies at high metallicities. We could wonder if gas-poor dwarf galaxies would show different, possibly lower, G/D than that observed in gas-rich dwarfs, thus possibly increasing the observed scatter at low metallicities. On the high-metallicity side, Smith et al. (2012b) showed that the 30 elliptical galaxies detected with Herschel in the Herschel Reference Survey (HRS, Boselli et al. 2010b) had a mean G/D of $\sim 120$, which is slightly lower than what we find for our elliptical and spiral galaxies at moderate metallicities. However, the dust masses were estimated via a modified blackbody model, thus we will not go deeper into any further comparison. Nonetheless, including more elliptical galaxies might also slightly increase the scatter at high metallicities.


Fig. 9.4. (Top row): G/D as a function of metallicity for the 2 values of $X_{C O}: X_{C O, M W}(a)$ and $X_{C O, Z}$ (b). The mean error for the data points is shown in grey on the right of the plots. The colours code the reliability of the point depending if the gas mass is uncertain (in blue), the dust mass is uncertain (in red) or if both are uncertain (in purple). The symbol traces the changes made in the HI and $\mathrm{H}_{2}$ masses (see text for details on the uncertainties and the changes on the gas masses). The dashed line represents the reference scaling of the G/D with metallicity (not fit to the data). The dotted and dash-dotted lines represent the best power-law and best broken power-law fits to the data. (Bottom row): Same as top row for $X_{C O, M W}(c)$ and $X_{C O, Z}(d)$, where the binned $\mathrm{G} / \mathrm{D}$ values (see text) have been added as pink filled circles. For clarity, the observed G/D values are now shown in grey. On the bottom panels the relative dispersion in each bins, in terms of standard deviation, is shown and the colours show the number of galaxies in each bin.

### 9.5.2 Empirical relations and scatter

To investigate the variation of the G/D with metallicity, we first fit a power law (dotted line in Fig. 9.4) through the observed $\mathrm{G} / \mathrm{D}$ values (excluding the limits): $\mathrm{G} / \mathrm{D} \propto(O / H)_{0}^{\alpha}$. The fit is performed with the IDL procedure MPFIT ${ }^{3}$ and is shown in Fig. 9.4. The fit is weighted by the individual errors bars of the G/D values and the number density of points to avoid being dominated by the more numerous high-metallicity galaxies. We get a slope for the power law of $\alpha_{0}=-1.6 \pm 0.3$ for $X_{C O, M W}$ and $\alpha_{0}=-2.0 \pm 0.3$ for $X_{C O, Z}$. In both cases, $\alpha_{0}$ is lower than -1 , which corresponds to the slope of the reference relation.

We also fit a broken power law (dash-dotted line in Fig. 9.4), with two slopes $\alpha_{L}$ and $\alpha_{H}$ to describe the low- and high-metallicity slopes respectively, and with a transition metallicity, $x_{t}$, between the two regimes. Several studies (e.g., James et al. 2002; Draine et al. 2007; Galliano et al. 2008; Leroy et al. 2011) have shown that the G/D was well represented by a power law with a slope of -1 at high metallicities and down to $12+\log (\mathrm{O} / \mathrm{H}) \sim 8.0-8.2$, and thus we fix $\alpha_{H}=-1$. This gives us a low-metallicity slope, $\alpha_{L}$, of $-3.1 \pm 1.8$ with a transition metallicity of $7.96 \pm 0.47$ for $X_{C O, M W}$ and $\alpha_{L}=-3.1 \pm 1.3$ and a transition around a metallicity of $8.10 \pm 0.43$ for $X_{C O, Z}$. The low-metallicity slopes, $\alpha_{L}$, are also for both cases lower than -1 . The parameters for the different empirical relations are given in Table 9.2.

Table 9.2. Parameters for the three empirical relations between the G/D and metallicity: power law (slope of -1 and free) and broken power law for the two $X_{C O}$ values.

| Parameters | $X_{C O, M W}$ case | $X_{C O, Z}$ case |
| :---: | :---: | :---: |
| Power law, slope fixed: $\mathrm{y}=\mathrm{a}+\left(\mathrm{x}_{\odot}-\mathrm{x}\right)$ ("reference" scaling) |  |  |
| $\mathrm{a}^{1,2}$ | 2.21 | 2.21 |
| average logarithmic distance ${ }^{3}$ [dex] | 0.07 | 0.27 |
| Power law, slope free: $\mathbf{y}=\mathbf{a}+\alpha_{0}\left(\mathbf{x}_{\odot}-x\right)$ |  |  |
| $\mathrm{a}^{1,2}$ | 2.21 | 2.21 |
| $\alpha_{0}$ | $1.62 \pm 0.34$ | $2.02 \pm 0.28$ |
| average logarithmic distance ${ }^{3}$ [dex] | -0.21 | -0.19 |
| Broken power law: |  |  |
| $\mathbf{y}=\mathbf{a}+\alpha_{H}\left(\mathbf{x}_{\odot}-x\right)$ for $\mathbf{x}>\mathbf{x}_{t}$ |  |  |
| $\mathbf{y}=\mathbf{b}+\alpha_{L}\left(\mathbf{x}_{\odot}-x\right)$ for $\mathbf{x} \leq \mathbf{x}_{t}$ |  |  |
| $\mathrm{a}^{1,2}$ | 2.21 | 2.21 |
| $\alpha_{H}$ | 1.00 | 1.00 |
| b | 0.68 | 0.96 |
| $\alpha_{L}$ | $3.08 \pm 1.76$ | $3.10 \pm 1.33$ |
| $\mathrm{x}_{t}$ | $7.96 \pm 0.47$ | $8.10 \pm 0.43$ |
| average logarithmic distance ${ }^{3}$ [dex] | -0.06 | 0.06 |

Notes. $\mathrm{y}=\log (\mathrm{G} / \mathrm{D}), \mathrm{x}=12+\log (\mathrm{O} / \mathrm{H})$ and $\mathrm{x}_{\odot}=8.69$.
${ }^{1}$ : Fixed parameters.
${ }^{2}$ : This correspond to the solar G/D: G/D $\odot=10^{a}=162$ (Zubko et al. 2004).
${ }^{3}$ : Derived for all of the individual galaxies, neglecting the upper/lower limits on the G/D.
If we let $\alpha_{H}$ free in the broken power-law fit, we get similar results within errors for $\alpha_{L}$ and $x_{t}$ in both $X_{C O}$ cases. We get $\alpha_{H}=-0.5 \pm 0.9$ and $\alpha_{H}=-1.6 \pm 0.6$, for $X_{C O, M W}$ and $X_{C O, Z}$ respectively, which is coherent with a slope of -1 within errors. Note that we have imposed here that our fits go through the solar G/D determined by Zubko et al. (2004). If we relax this condition (i.e., do not

[^48]fix our "a" parameter in Table 9.2), we get values of the solar G/D ranging from $(G / D)_{\odot}=90$ to 240 within $\sim 60 \%$ of the value from Zubko et al. (2004).

Now we consider the deviation from each relation by looking at the logarithmic distance from the observed G/D values and the G/D values predicted by each of the three relations presented in Table 9.2. This is a way to look at the residuals from the two fits and the reference scaling, even though we did not actually fit the reference trend to the $G / D$ values. These residuals are shown in Fig. 9.5. Average residuals in each metallicity bin defined previously are also computed. For a given point, the best relation is the one giving the residual closest to zero. From Fig. 9.5 we have another confirmation that a reference scaling of the $G / D$ with metallicity does not provide reliable estimates of the G/D at low metallicities $12+\log (\mathrm{O} / \mathrm{H}) \lesssim 8.4$. We also note that, for the average residuals, the broken power law gives the residuals that are the closest to zero for nearly all the metallicity bins in both $X_{C O, M W}$ and $X_{C O, Z}$ cases, and that this corresponds to a predicted G/D uncertain to a factor of 1.6.

Even though $30 \%$ of our sample have metallicities below $1 / 5 Z_{\odot}$, only seven galaxies have $12+\log (\mathrm{O} / \mathrm{H}) \leq 7.5$ with two of them not detected in Hi (SBS1159+545 and Tol1214-277). The remaining five galaxies (IZw18, HS0822+3542, SBS0335-052, SBS1415 +437 and UGC4483) are important constraints for the broken power-law fit. These five galaxies all present broad dust SEDs peaking at very short wavelengths ( $\sim 40 \mu \mathrm{~m}$, and $\sim 70 \mu \mathrm{~m}$ for HS0822+3542), indicating overall warmer dust with a wide range of dust grain temperatures, and subsequently very low dust masses; hence their high G/D. This peculiar SED shape had already been noted by Rémy-Ruyer et al. (2013).


Fig. 9.5. Residuals (i.e., logarithmic distance) between the observed (and detected) G/D and predicted G/D for the three relations for $X_{C O, M W}$ (left) and $X_{C O, Z}$ (right): reference scaling of the G/D with metallicity (crosses), the best power-law fit (filled circles) and the best broken power-law fit (triangles). These residuals are shown in grey for the individual galaxies and in colour for the average residuals in each metallicity bin defined in Section 9.5.1. The colours show the number of galaxies in each bin. The mean residual for all of the observed G/D values is shown by the dashed (reference scaling), dotted (power-law fit) and dash-dotted (broken power-law fit) lines for the three relations and are reported in Table 9.2.

Using Herschel data and a semi-empirical SED model, Sandstrom et al. (2013) looked at the G/D in a sub-sample of 26 KINGFISH galaxies, mostly spirals. They simultaneously derive $X_{C O}$ and G/D for their sample, taking advantage of the high spatial resolution of the KINGFISH gas and dust data. They found that the G/D for these galaxies follows the reference trend with the metallicity and shows small scatter. Their metallicity range is from $12+\log (\mathrm{O} / \mathrm{H}) \sim 8.1$ to 8.8 and thus these results are in agreement with our findings. Moreover the small scatter (less than a factor
of 2) observed by Sandstrom et al. (2013) can be due to the fact that they are probing very similar environments. In our case we have a wide variety of morphological types represented in our sample, that results in a larger scatter (a factor of $\sim 5$ and 3 for $X_{C O, M W}$ and $X_{C O, Z}$ respectively for this metallicity range).

### 9.5.3 Discussion

In the previous section, we have shown that the reference scaling relation between metallicity and G/D derived for metallicities above $12+\log (\mathrm{O} / \mathrm{H}) \sim 8.0$ does not apply to objects with lower metallicity. We empirically derived a new scaling relation better described by a broken power law with a transition metallicity around $12+\log (\mathrm{O} / \mathrm{H}) \sim 8.0$, which confirms the importance of this value in low-metallicity dwarf galaxies. As mentioned in Section 9.1, this reference scaling relation arises from the hypothesis that the dust formation timescale and the dust destruction timescale behave similarly with time. Thus a possible interpretation of our results would be that the balance between formation and destruction of dust grains is altered at low metallicity, resulting in the observed steeper trend. Dwarf galaxies are subject to an overall harder ISRF than more metal-rich environments (Madden et al. 2006). The harder UV photons travel deeper into the ISM and photoprocess dust in much deeper regions in the clouds limiting the accretion and coagulation of the grains. The hard ISRF also affects the dust survivability in such extreme environments, especially carbonaceous dust: the dust destruction by hard UV photons is enhanced in low-metallicity galaxies for small carbon grains (e.g., Pilleri et al. 2012; Bocchio et al. 2012, 2013). In dwarf galaxies, dust destruction by SN shocks is enhanced too compared to larger scale galaxies, as most of the ISM can be affected by the shock due to the small physical size of the dwarfs and due to the globally lower density of the ISM.

In the following paragraphs we discuss the impact of several assumptions made to estimate the G/D on our results: the dust composition, the choice of the radiation field for the dust modelling, and the potential presence of a submm excess in some of our dwarf galaxies.

Dust composition - Galliano et al. (2011) demonstrated that a more emissive dust grain composition compared to that of the Galaxy, is more consistent for the low-metallicity Large Magellanic Cloud (LMC). This result has been confirmed by Galametz et al. (2013) in a star-forming complex of the LMC with an updated version of the SPIRE calibration ${ }^{4}$, and by our study of the submm excess in Chapter 8. Changing accordingly the dust composition in our low-metallicity galaxies would give lower dust masses (by a factor of $\sim$ two) with more emissive dust grains and would also increase the G/D by the same factor, increasing the discrepancy at low metallicities between the observed G/D and the predicted G/D from the reference scaling relation.

Radiation field - In the dust modelling, we use an ISRF with the spectral shape of the Galactic ISRF for all of our galaxies for consistency. However, the ISRF in low-metallicity dwarf galaxies is harder, so we could wonder if this spectral shape is appropriate for the modelling of dwarf galaxies. As already mentioned in Chapter 7, the shape of the radiation field determines the emission of out-of-equilibrium small grains. Increasing the hardness of the radiation field increases the maximum temperature the small grains can reach when they undergo stochastic heating. However, these very small grains only have a minor contribution to the total dust mass, and thus the assumed shape of the ISRF does not bias our estimation of the total dust mass for dwarf galaxies.

[^49]Submm excess - A submm excess has been observed in the past in several low-metallicity galaxies that current dust SED models are unable to fully explain (Galliano et al. 2003, 2005; Dumke et al. 2004; Bendo et al. 2006; Zhu et al. 2009; Galametz et al. 2009; Bot et al. 2010; Grossi et al. 2010; Galametz et al. 2011). We confirmed the presence of such an excess in several DGS and KINGFISH galaxies (see Chapter 8). Several hypothesis have been made to explain this excess among which the addition of a VCD component in which most of the dust mass should reside. This VCD component would be in the form of very dense clumps in the ISM (Galliano et al. 2003, 2005). Taking this additional VCD component into account in the DGS and KINGFISH galaxies presenting a submm excess can result in a drastic increase of the dust mass and thus in a lower G/D. However, Galliano et al. (2011) showed that for a strip of the LMC, the submm excess is more important in the diffuse regions, possibly in contradiction with the hypothesis of very cold dust in dense clumps. Other studies have suggested an enhanced fraction of very small grains with high emissivity (Lisenfeld et al. 2002; Dumke et al. 2004; Bendo et al. 2006; Zhu et al. 2009), "spinning" dust emission (Ysard \& Verstraete 2010) or emission from magnetic nano-particles (Draine \& Hensley 2012) to explain the submm excess. Meny et al. (2007) proposed variations of the optical properties of the dust with the temperature which results in an enhanced emission of the dust at submm $/ \mathrm{mm}$ wavelengths. We saw in Chapter 8 that a dust composition using amorphous carbon grains could explain the excess for some moderately metal-poor galaxies. Using dust masses estimated with amorphous carbon grains for the 37 galaxies presenting a submm excess would result in an increase of a factor of $\sim 2$ for the G/D of these galaxies, which is within the scatter of the G/D vs metallicity relation, and hence would not change our general conclusions.

### 9.6 Chemical evolution models

Chemical evolution models, under certain assumptions, can predict a possible evolution of the G/D as metallicity varies in a given galaxy. For example, in the disk of our Galaxy, chemical evolution models predict this "reference" scaling of the G/D with the metallicity (Dwek 1998). We consider three different models here, from Galliano et al. (2008), Asano et al. (2013) and Zhukovska (2014) to interpret our data. However, we have to keep in mind during this comparison that, since we do not know the ages of these galaxies and that the same metallicity can be reached at very different time by different galaxies, our sample cannot be considered as the evolution (snapshots) of one single galaxy.

### 9.6.1 A simple model to begin with

Galliano et al. (2008) developed a one-zone single-phase chemical evolution model, based on the model by Dwek (1998). They consider a closed-box model where the evolution of the dust content is regulated by balancing dust production by stars and dust destruction by star formation and SN blast waves. They assume the full condensation of the elements injected by Type II supernovae (SNII) into dust and instantaneous mixing of the elements in the ISM. The model is shown on Fig. 9.6 for various SN destruction efficiencies as the dark grey zone.

Two things can be noticed from Fig. 9.6. First, the model is consistent with the observed G/D at high metallicities within the scatter, and down to metallicities $\sim 0.5 Z_{\odot}$. Second, the model does not work at low metallicities and systematically underestimates the G/D. This has already been noted by Galliano et al. (2008) for their test sample of galaxies and was attributed to the very crude assumptions made in the modelling, especially the instantaneous mixing of the SNII elements in the ISM. Another strong simplifying assumption made by Galliano et al. (2008) is that they did not take into account dust growth in the ISM as they assume full condensation of the grains. In


Fig. 9.6. G/D as a function of metallicity for the 2 values of $X_{C O}: X_{C O, M W}(t o p)$ and $X_{C O, Z}$ (bottom) with the chemical evolution model of Galliano et al. (2008). The colours and symbols are the same as for Fig.9.4. The dark grey stripes show the range of values from the Galliano et al. (2008) chemical evolution model. The black dashed line represents the reference scaling of the $\mathrm{G} / \mathrm{D}$ with metallicity (not fit to the data). The black dotted and dash-dotted lines represent the best power-law and best broken power-law fits to the data.
the Galaxy, the typical timescale for dust formation by stars has been shown to be larger than the typical timescale for dust destruction (Jones \& Tielens 1994; Jones et al. 1996). Because we still observe dust in the ISM, we need to reach equilibrium between formation and destruction of the dust grains, either via high SN yields or dust growth processes in the ISM. In the following sections we thus look at models including dust growth in the ISM.

### 9.6.2 Including dust growth in the ISM

Asano et al. (2013) propose a chemical evolution model, based on models from Hirashita (1999) and Inoue (2011), taking into account the evolution of the metal content in the dust phase in addition to the evolution of the total amount of metals. The dust formation is regulated by AGB stars, SNII and dust growth, via accretion, in the ISM. The dust is destroyed by SN shocks. Inflows and outflows are not considered (closed-box model) and the total mass of the galaxy is constant and set to $10^{10} M_{\odot}$. Metallicity and age dependence of the various dust formation processes are taken into account. Asano et al. (2013) show that dust growth in the ISM becomes the main driver of the dust mass evolution, compared to the dust formation from metals produced and ejected into the ISM by stars, when the metallicity of the galaxy exceeds a certain "critical" metallicity. This critical metallicity increases with decreasing star-formation timescale. Asano et al. (2013) show that dust growth via accretion processes in the ISM is regulated by this critical metallicity over a large range of star-formation timescales (for $\tau_{\mathrm{SF}}{ }^{5}=0.5-5-50 \mathrm{Gyr}$ ). After reaching this critical metallicity the dust mass increases more rapidly, boosted by dust growth processes, before saturating when all of the metals available for dust formation are locked up in dust. The metallicity at which this saturation occurs thus also depends on the critical metallicity, which in turn depends on the star-formation history of the galaxy.

Figure 9.7 shows the models of Asano et al. (2013) (for $\tau_{\mathrm{SF}}=0.5-5-50 \mathrm{Gyr}$ ) overlaid on the observed G/D values. The models were originally on an arbitrary scale and they are thus normalised at the $(G / D)_{\odot}$ value. We assume an error on this value of $\sim 60 \%$, from the range of values determined from the fits in Section 9.5.2, to have a tolerance range around the model (shown by the shaded grey area on Fig. 9.7). The three models show similar evolution with metallicity and indeed are homologous to each other when normalised by their respective critical metallicities (see Fig. 3 of Asano et al. 2013). We clearly see the influence of the critical metallicity on the dust mass evolution: at low metallicities the range of possible G/D values (illustrated by the grey area on Fig. 9.7) becomes wider around $12+\log (\mathrm{O} / \mathrm{H}) \sim 7.2-7.3$ before narrowing down around $12+\log (\mathrm{O} / \mathrm{H}) \sim 8.6$. This broadening is due to the fact that in this range of metallicities, galaxies with high star-formation timescales have already reached their critical metallicity and have a rapidly increasing dust mass (and thus a low G/D at a given metallicity), compared to galaxies with lower star-formation timescales which have not yet reached this critical metallicity and with a dust mass still regulated by stars (thus with a higher G/D at the same metallicity). Galaxies with high starformation timescales then reach saturation at moderate metallicities as they started their "active dust growth" phase at a lower critical metallicity (i.e., earlier in their evolution), while, at the same metallicity, galaxies with low star-formation timescales are still in the "active dust growth" phase. Then when these galaxies also reach saturation, because the dust growth in the ISM becomes ineffective, the range of possible G/D values narrows down.

From Fig. 9.7 we see that the models from Asano et al. (2013) are consistent with the G/D from both $X_{C O}$ values. We note that below $12+\log (\mathrm{O} / \mathrm{H}) \sim 7.5$, even the $\tau_{\mathrm{SF}}=50 \mathrm{Gyr}$ model does not agree anymore with the reference scaling of the G/D with metallicity (and below $12+\log (\mathrm{O} / \mathrm{H}) \sim$ 8.0 for $\tau_{\mathrm{SF}}=5 \mathrm{Gyr}$ ). The other two empirical relations (our best power-law and broken power-law

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Fig. 9.7. G/D as a function of metallicity for the 2 values of $X_{C O}: X_{C O, M W}(t o p)$ and $X_{C O, Z}$ (bottom) with the chemical evolution model of Asano et al. (2013). The symbols are the same as for Fig. 9.4. The colours delineate ranges in star-formation timescales $\tau_{\mathrm{SF}}$. The model from Asano et al. (2013) is overlaid on the points for various $\tau_{\mathrm{SF}}=0.5$ (red), 5 (blue), 50 (purple) Gyr. The black dashed line represents the reference scaling of the G/D with metallicity (not fit to the data). The black dotted and dash-dotted lines represent the best power-law and best broken power-law fits to the data.
fits) are consistent with the models of Asano et al. (2013) within the considered metallicity range: from $12+\log (\mathrm{O} / \mathrm{H}) \sim 7.0$ to 9.1. Comparing with the shape of the Asano et al. (2013) models, our best broken power-law fit may overestimate the G/D for $12+\log (\mathrm{O} / \mathrm{H}) \leq 7.0$. However, Izotov et al. (2012) recently suggested that there seems to be a metallicity floor around $12+\log (\mathrm{O} / \mathrm{H}) \sim 6.9$, below which no galaxies are found in the local Universe, as already proposed by Kunth \& Sargent (1986). Thus our metallicity range is close to being the largest achievable in the local Universe as far as low metallicities are concerned.

The galaxies from the DGS, KINGFISH and G11 samples are colour coded in Fig. 9.7 by an approximation of their star-formation timescale $\tau_{\mathrm{SF}}$, estimated from $\tau_{\mathrm{SF}}=\left(M_{g a s}+M_{\text {dust }}\right) / \mathrm{SFR}$, where the SFR have been estimated from $\mathrm{L}_{T I R}$ (see Chapter 7). The $\tau_{\mathrm{SF}}$ values are roughly consistent with the models from Asano et al. (2013). The median value of $\tau_{\mathrm{SF}}$ is $\sim 3.0\left(X_{C O, M W}\right)$ and $5.5\left(X_{C O, Z}\right)$ Gyr, but with a large dispersion of $\sim 20 \mathrm{Gyr}$ around this value. As the models from Asano et al. (2013) encompass most of the observed G/D values, the dispersion seen in the G/D values can be due to the wide range of star-formation timescales in the considered galaxies. This is consistent with the large dispersion in the approximated star-formation timescales in our sample.

In Asano et al. (2013), the star formation is assumed to be continuous over the star-formation timescale. However, star-formation histories of many dwarf galaxies derived from colour-magnitude diagrams reconstruction show distinct episodes of star formation separated by more quiescent phases (e.g., Tolstoy et al. 2009and references therein). For example, Legrand et al. (2000) suggested for IZw18 a star-formation history made of bursts of star formation in between more quiescent phases, following the suggestion of Searle \& Sargent (1972). Episodic star-formation histories have also been suggested in Nbody/Smoothed Particle Hydrodynamics simulations of dwarf galaxy evolution (e.g., Valcke et al. 2008; Revaz \& Jablonka 2012). As we saw that the scatter in Fig. 9.7 seems to be due to the range of star-formation timescales probed by our sample, and assuming a continuous star formation, we thus need to consider the influence of the continuous vs. episodic star-formation modes.

### 9.6.3 Episodic versus continuous star formation

In the following we compare the observationally derived G/D with results of dust evolution models in dwarf galaxies with episodic star-formation history from Zhukovska (2014). These models were originally introduced to study the lifecycle of dust species from different origins in the Solar neighbourhood (Zhukovska et al. 2008). The model of Zhukovska (2014) is based on the Zhukovska et al. (2008) model that has been adapted to treat dwarf galaxies, specifically by considering episodic star formation. In Zhukovska (2014), the equations describing the evolution of the galaxy are now normalised to the total galactic masses $M_{\text {tot }}$ (instead of surface densities) because dwarf galaxies are smaller in size and thus assumed to have a well mixed ISM. As in Zhukovska et al. (2008), the modelled dwarf galaxy is formed by gas infall starting from $M_{\text {tot }}=0$ and reach its total mass $M_{\text {tot }}$ on the infall timescale. The assumed value of the infall timescale in Zhukovska (2014) is set to a much shorter value for dwarf galaxies than for the Solar neighbourhood. Since the G/D is the ratio of the gas and dust masses, it does not depend on the normalisation by the total mass $M_{\text {tot }}$, and is determined by the star-formation history and infall timescale. We refer the reader to Zhukovska (2014), for more details on the modelling.

Similarly to the models from Asano et al. (2013), Zhukovska (2014) include dust formation in AGB stars, SN II and dust growth by mantle accretion in the ISM. The main difference between these models is in the treatment of dust growth. Zhukovska (2014) assume a two-phase ISM consisting of clouds and an intercloud medium, where clouds are characterised by temperature, density, mass fraction, and lifetime. Dust growth by accretion in their model takes place only in the dense gas and also critically depends on the metallicity (see Zhukovska 2008).


Fig. 9.8. G/D as a function of metallicity for the 2 values of $X_{C O}: X_{C O, M W}(t o p)$ and $X_{C O, Z}($ bottom $)$ with the chemical evolution model of Zhukovska (2014). The symbols are the same as for Fig. 9.4. The model from Zhukovska (2014) is shown for various star-formation histories: episodic with 6 bursts of 50 Myr and star-formation timescale of $\tau_{\mathrm{SF}}=2 \mathrm{Gyr}$ (orange), episodic with 6 bursts of 500 Myr and $\tau_{\mathrm{SF}}=2 \mathrm{Gyr}$ (green), episodic with 6 bursts of 500 Myr and $\tau_{\mathrm{SF}}=0.2 \mathrm{Gyr}$ (more intense star formation, brown) and continuous with a star-formation timescale $\tau_{\mathrm{SF}}=10 \mathrm{Gyr}$ (cyan dash- 3 dots line). The black dashed line represents the reference scaling of the G/D with metallicity (not fit to the data). The black dotted and dash-dotted lines represent the best power-law and best broken power-law fits to the data.

In this paper we consider three models from Zhukovska (2014), which differ only in duration and intensity of the star formation bursts. All models consider six bursts of star formation starting at instants $t=0.5,1,2,5,7$, and 11 Gyr. In the first and the second model, the burst duration is 50 Myr and 500 Myr , respectively, and the $\tau_{\mathrm{SF}}$ during bursts is 2 Gyr . The first model is typical of a low-metallicity dwarf galaxy. In the third model, the burst duration is 500 Myr but the value of $\tau_{\mathrm{SF}}$ is much shorter, 0.2 Gyr. During the quiescence phases $\tau_{\mathrm{SF}}$ is set to be 200 Gyr . We also consider a model with continuous star formation on a 10 Gyr timescale, for comparison. In all models, the infall timescale is 0.3 Gyr and there are no galactic outflows. The initial metallicity $Z$ of the infalling gas is set to be $10^{-4}$ with SNII like enhanced $[\alpha / \mathrm{Fe}]$ ratio.

The models from Zhukovska (2014) are presented in Fig. 9.8 and reproduce the broadening of the observed G/D values at low metallicities $(12+\log (\mathrm{O} / \mathrm{H}) \lesssim 8.3)$, and also converge around $12+\log (\mathrm{O} / \mathrm{H}) \sim 7.2$, similar to the models of continuous star formation of Asano et al. (2013). Note how the star-formation history impacts the shape of the modelled G/D: the most extreme G/D values are obtained by the three models with episodic bursts of star formation. For the model with more intense star formation (brown curve on Fig. 9.8), $12+\log (\mathrm{O} / \mathrm{H})=8.6$ is reached during the first burst, and very high values of the G/D are quickly reached, up to two orders of magnitude above the reference scaling relation at moderate metallicities $(12+\log (\mathrm{O} / \mathrm{H}) \sim 8.2-8.3)$. It also presents an interesting scatter of G/D values near $12+\log (\mathrm{O} / \mathrm{H})=9.0$ that is due to dust destruction during the SF bursts, and consistent with the scatter predicted by the Galliano et al. (2008) model (see Fig. 9.9). The fact that the low-metallicity slope of the broken power law is consistent with the continuous star formation model at low metallicities for the $X_{C O, Z}$ case indicates that this broken power law can provide a fairly good empirical way of estimating the G/D for a given metallicity.

### 9.6.4 Explaining the observed scatter in G/D values

Figure 9.9 shows the three models overlaid on the observed G/D values. The models from Asano et al. (2013) and Zhukovska (2014) provide trends that are consistent with each other and with the data and its scatter. More dust observations of extremely low-metallicity galaxies with $12+\log (\mathrm{O} / \mathrm{H})$ $<7.5$ are nonetheless needed to confirm this agreement between the models and the data at very low metallicities. The model of Galliano et al. (2008) fails to reproduce the observed G/D at low metallicities, but provides a good complement to explain the scatter seen at high metallicities, consistent with the predictions of the third bursty model by Zhukovska (2014). We thus conclude that the observed scatter at low metallicities in the G/D values is due to the wide variety of environments we are probing, and especially to the different star-formation histories. The observed scatter at higher metallicity seems to be due to different timescales for dust destruction by SN blast waves in the different environments and to the efficiency of dust shattering in the ISM.

We investigated here two different parameters to explain the scatter in the G/D values: starformation histories and efficiency of dust destruction, but other processes could also give rise to the observed scatter. In our dust modelling we allow the mass fraction of small grains compared to big grains to vary from galaxy to galaxy (controlled by the $f_{v s g}$ parameter), to account for potential variations in the grain size distribution. This had already been done in one low-metallicity galaxy by Lisenfeld et al. (2002). On the theoretical side, Hirashita \& Kuo (2011) showed that the dust grain size distribution can have an important impact on the dust growth process in the ISM by regulating the grain growth rate. They also showed that the critical metallicity mentioned in Section 9.6.2, for which grain growth becomes dominant, also depends on the grain size distribution. Additionally the grain size distribution varies as the galaxy evolves and this evolution is controlled by different dust formation processes at different ages (Asano et al. 2013). Thus the observed scatter can also be due to variations of the grain size distribution between the galaxies, the effect of which can be related to the star-formation history.

Second, as discussed in Section 9.5 .3 we use a dust model with the dust composition and optical


Fig. 9.9. G/D as a function of metallicity for the 2 values of $X_{C O}: X_{C O, M W}(t o p)$ and $X_{C O, Z}$ (bottom) with the three chemical evolution model considered in Section 9.6. The symbols are the same as for Fig. 9.4. The model ranges from Galliano et al. (2008) are delineated by the dark grey stripe. The models from Asano et al. (2013) are shown with the red, blue and purple lines. The models from Zhukovska (2014) are shown with the orange, green brown solid lines and cyan dash-3 dots line. The black dashed line represents the reference scaling of the G/D with metallicity (not fit to the data).
properties representative of dust in the Milky Way for all of the galaxies. Another explanation for the scatter seen at all metallicities could be that the dust composition in fact varies between the galaxies, leading to important variations in the emissivity of the dust grains (Jones 2012). This would then imply dust masses relatively similar at a given metallicity but large variations in the emissivity properties of dust from one galaxy to another. With our fixed emissivity (due to our fixed dust composition in our dust model) this effect would be seen through the variations in the derived dust masses (and then in the G/D values) seen at a given metallicity, thus giving rise to the observed scatter.

Another aspect not taken into account in the chemical models we use here is that the mass of the galaxy plays an important role in its chemical evolution. According to the chemical downsizing scenario, galaxies with different masses have different star-formation efficiencies (e.g., Brooks et al. 2007). The massive galaxies form stars before low-mass galaxies in the history of the Universe (Cowie et al. 1996). The exact reasons are not well known, some models would argue this is a feedback effect, with metals being lost in outflows in less massive galaxies (Frye et al. 2002), and others would argue that the local density and the ISM pressure is responsible (e.g., Cen \& Ostriker 1999): if the ISM density is systematically lower in a dwarf galaxy, for instance, then the starformation activity will be low on average, impacting the G/D value. A massive galaxy will not show the same behaviour, and this difference will also introduce some scatter in the observed G/D at a given metallicity.

Additionally external processes such as outflows or interactions and mergers, not considered here either in our models, can also be responsible for the scatter in the G/D values.

### 9.6.5 Implications for the observed G/D in galaxies

We saw in Sections 9.6.2 and 9.6.3 that the broken power-law relation in the $X_{C O, Z}$ case is consistent with the predictions from both chemical evolution models. Even though there are only five galaxies with $12+\log (\mathrm{O} / \mathrm{H}) \sim 7.5$ to constrain the observed G/D at extremely low metallicities, this broken power-law relation is the best empirical estimate of the observed G/D for local galaxies we have at our disposal so far. Thus, we advise to use this empirical prescription to estimate the G/D based on a metallicity value for local galaxies, keeping in mind the large scatter and uncertainties on the broken power-law parameters: the estimated G/D would be accurate to a factor of $\sim 1.6$. Note also that this empirical relation has been derived for metallicities estimated with a strong emission line method and the calibration from Pilyugin \& Thuan (2005). Thus the estimation of the G/D with this empirical relation should be done from a metallicity derived with the same method. Additionally, more observations of the dust in extremely low-metallicity galaxies are needed in order to place more constraints on this empirical relation.

Chapter 9. Gas-to-Dust mass ratios as a function of metallicity
Table 9.3. Gas masses for the DGS, KINGFISH and G11 samples.

|  | $M_{H I}\left[M_{\odot}\right]$ |  |  | $M_{H_{2}}\left[M_{\odot}\right]$ |  |  | $\mathbf{M}_{\text {gas }}\left[M_{\odot}\right]$ |  |  |  |  |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Name | $M_{H I}$ (Ref) | Unc (\%) | HIarea ( $\operatorname{arcmin}^{2}$ ) | $12+\log (\mathrm{O} / \mathrm{H})(\mathrm{Ref})$ | $M_{H_{2}, M W}$ (Ref) | $M_{H_{2}, Z}$ | $\mathrm{M}_{\text {gas }, M W}$ | Unc (\%) | $\mathrm{M}_{\text {gas }, Z}$ | Unc (\%) | $\mu_{\text {gal }}$ |
| DGS |  |  |  |  |  |  |  |  |  |  |  |
| Harol1 | $5.01 \times 10^{8 a, e}$ (1) | 50.00 | - | 8.36 (1) | $2.51 \times 10^{8}$ (1) | $1.17 \times 10^{9}$ | $1.74 \times 10^{9}$ | 19.94 | $3.01 \times 10^{9}$ | 11.50 | 1.38 |
| Haro2 | $3.77 \times 10^{8 c}(2)$ | 16.08 | 1.6 | 8.23 (1) | $1.26 \times 10^{8}$ (2) | $1.03 \times 10^{9}$ | $6.93 \times 10^{8}$ | 12.06 | $1.94 \times 10^{9}$ | 4.31 | 1.38 |
| Haro3 | $1.13 \times 10^{9 a}$ (3) | 2.13 | - | 8.28 (1) | $3.02 \times 10^{7}$ (3) | $2.01 \times 10^{8}$ | $1.60 \times 10^{9}$ | 2.08 | $1.84 \times 10^{9}$ | 1.81 | 1.38 |
| He2-10 | $3.10 \times 10^{8 c}$ (4) | 6.45 | 3.2 | 8.43 (1) | $9.55 \times 10^{7}$ (4) | $3.21 \times 10^{8}$ | $5.61 \times 10^{8}$ | 4.93 | $8.73 \times 10^{8}$ | 3.17 | 1.39 |
| HS0017 + 1055 | $\leq 1.74 \times 10^{8 a}$ (5) | - | - | 7.63 (1) | $\leq 2.15 \times 10^{6 d}$ (-) | $\leq 1.18 \times 10^{8}$ | $\leq 2.42 \times 10^{8}$ | - | $\leq 4.01 \times 10^{8}$ | - | 1.37 |
| HS0052+2536 | $\leq 4.82 \times 10^{10 a}(6)$ | - | - | 8.04 (1) | $\leq 5.94 \times 10^{8 d}(-)$ | $\leq 3.26 \times 10^{10}$ | $\leq 6.72 \times 10^{10}$ | - | $\leq 1.11 \times 10^{11}$ | - | 1.38 |
| HS0822 +3542 | $5.75 \times 10^{7 b}(7)$ | 22.22 | 0.6 | 7.32 (1) | $9.48 \times 10^{5 d}(-)$ | $5.21 \times 10^{7}$ | $8.02 \times 10^{7}$ | 21.86 | $1.50 \times 10^{8}$ | 11.66 | 1.37 |
| HS1222+3741 | - | - | - | 7.79 (1) | - | - | - | - | - | - | 1.37 |
| HS1236+3937 | - | - | - | 7.72 (1) | - | - | - | - | - | - | 1.37 |
| HS1304+3529 | - | - | - | 7.93 (1) | - | - | - | - | - | - | 1.38 |
| HS1319+3224 | - | - | - | 7.81 (1) | - | - | - | - | - | - | 1.37 |
| HS1330 + 3651 | - | - | - | 7.98 (1) | ${ }^{-}$ | - | - | - | - | - | 1.38 |
| HS1442+4250 | $3.10 \times 10^{8 a}$ (5) | 2.27 | - | 7.60 (1) | $3.82 \times 10^{6 d}(-)$ | $2.10 \times 10^{8}$ | $4.30 \times 10^{8}$ | 2.24 | $7.13 \times 10^{8}$ | 1.35 | 1.37 |
| HS2352+2733 | $\leq 2.30 \times 10^{10 a}$ (6) | - | - | 8.40 (1) | - | - | - | - | - | - | 1.38 |
| IZw18 | $1.00 \times 10^{8 c 2}$ (8) | 16.08 | - | 7.14 (1) | $2.71 \times 10^{6 d}$ (5) | $1.49 \times 10^{8}$ | $1.41 \times 10^{8}$ | 15.66 | $3.41 \times 10^{8}$ | 6.46 | 1.37 |
| IC10 | $4.41 \times 10^{7 c 2}(9)$ | 4.52 | - | 8.17 (1) | $4.84 \times 10^{6}$ (6) | $5.27 \times 10^{7}$ | $6.74 \times 10^{7}$ | 4.07 | $1.33 \times 10^{8}$ | 2.06 | 1.38 |
| IIZw40 | $5.66 \times 10^{8 c}$ (10) | 16.08 | 0.3 | 8.23 (1) | $2.83 \times 10^{7}$ (7) | $2.41 \times 10^{8}$ | $8.20 \times 10^{8}$ | 15.32 | $1.11 \times 10^{9}$ | 11.29 | 1.38 |
| Mrk1089 | $1.47 \times 10^{10 b}$ (11) | 7.91 | 26.4 | 8.10 (1) | $2.84 \times 10^{8}(1)$ | $4.31 \times 10^{9}$ | $2.07 \times 10^{10}$ | 7.76 | $2.62 \times 10^{10}$ | 6.12 | 1.38 |
| Mrk1450 | $4.30 \times 10^{7 c}$ (12) | 16.08 | 0.1 | 7.84 (1) | $5.30 \times 10^{5 d}(-)$ | $2.91 \times 10^{7}$ | $5.98 \times 10^{7}$ | 15.89 | $9.90 \times 10^{7}$ | 9.59 | 1.37 |
| Mrk153 | $\leq 6.93 \times 10^{8 a}$ (13) | - | - | 7.86 (1) | $\leq 8.39 \times 10^{6 d}(-)$ | $\leq 4.61 \times 10^{8}$ | $\leq 9.55 \times 10^{8}$ | - | $\leq 1.58 \times 10^{9}$ | - | 1.37 |
| Mrk209 | $2.76 \times 10^{7 b}$ (13) | 6.58 | 12.2 | 7.74 (1) | $3.09 \times 10^{6}$ (8) | $2.50 \times 10^{8}$ | $4.21 \times 10^{7}$ | 5.92 | $3.81 \times 10^{8}$ | 0.66 | 1.37 |
| Mrk930 | $3.19 \times 10^{9 c}(14)$ | 11.06 | 0.1 | 8.03 (1) | $3.07 \times 10^{7}$ (1) | $6.57 \times 10^{8}$ | $4.43 \times 10^{9}$ | 10.96 | $5.30 \times 10^{9}$ | 9.17 | 1.38 |
| NGC1140 | $3.50 \times 10^{9 a}$ (15) | 26.71 | - | 8.38 (1) | $1.91 \times 10^{7}(9)$ | $8.02 \times 10^{7}$ | $4.87 \times 10^{9}$ | 26.57 | $4.95 \times 10^{9}$ | 26.11 | 1.38 |
| NGC1569 | $1.77 \times 10^{8 b}(16)$ | 16.08 | 24.0 | 8.02 (1) | $7.08 \times 10^{5}(10)$ | $1.54 \times 10^{7}$ | $2.45 \times 10^{8}$ | 16.02 | $2.65 \times 10^{8}$ | 14.80 | 1.38 |
| NGC1705 | $7.67 \times 10^{7 b}(17)$ | 12.50 | 13.0 | 8.27 (1) | -08×105 | - | $\geq 1.06 \times 10^{8}$ | - | $\geq 1.06 \times 10^{8}$ | - | 1.38 |
| NGC2366 | $2.93 \times 10^{8 b}$ (18) | 6.76 | 420.0 | 7.70 (1) | $9.12 \times 10^{6 d}$ (11) | $5.01 \times 10^{8}$ | $4.15 \times 10^{8}$ | 6.55 | $1.09 \times 10^{9}$ | 2.49 | 1.37 |
| NGC4214 | $3.76 \times 10^{8 b}$ (16) | 16.08 | 156.0 | 8.26 (1) | $1.62 \times 10^{6}(12)$ | $1.16 \times 10^{7}$ | $5.22 \times 10^{8}$ | 16.02 | $5.36 \times 10^{8}$ | 15.61 | 1.38 |
| NGC4449 | $9.55 \times 10^{8 b}$ (16) | 16.08 | 144.0 | 8.20 (1) | $2.67 \times 10^{7}(13)$ | $2.50 \times 10^{8}$ | $1.35 \times 10^{9}$ | 15.65 | $1.66 \times 10^{9}$ | 12.75 | 1.38 |
| NGC4861 | $4.12 \times 10^{8 b}$ (19) | 6.27 | 34.3 | 7.89 (1) | $5.90 \times 10^{6 d}(1)$ | $3.24 \times 10^{8}$ | $5.75 \times 10^{8}$ | 6.18 | $1.01 \times 10^{9}$ | 3.51 | 1.37 |
| NGC5253 | $1.06 \times 10^{8 b}(20)$ | 5.66 | 77.0 | 8.25 (1) | $3.02 \times 10^{6}(14)$ | $2.26 \times 10^{7}$ | $1.51 \times 10^{8}$ | 5.50 | $1.78 \times 10^{8}$ | 4.67 | 1.38 |
| NGC625 | $1.10 \times 10^{8 c}(21)$ | 18.18 | 20.0 | 8.22 (1) | $4.36 \times 10^{6}(1)$ | $3.72 \times 10^{7}$ | $1.58 \times 10^{8}$ | 17.49 | $2.03 \times 10^{8}$ | 13.59 | 1.38 |
| NGC6822 | $1.04 \times 10^{8 b}(22)$ | 10.45 | 676.0 | 7.96 (1) | $3.58 \times 10^{5}(15)$ | $1.05 \times 10^{7}$ | $1.43 \times 10^{8}$ | 10.41 | $1.57 \times 10^{8}$ | 9.48 | 1.38 |
| Pox186 | $\leq 2.00 \times 10^{6 a, e}(23)$ | - | - | 7.70 (1) | $\leq 2.47 \times 10^{4 d}(-)$ | $\leq 1.35 \times 10^{6}$ | $\leq 4.74 \times 10^{6}$ | - | $\leq 6.57 \times 10^{6}$ | - | 1.37 |
| SBS0335-052 | $4.37 \times 10^{8 b}(24)$ | 10.27 | 0.2 | 7.25 (1) | $5.67 \times 10^{6 d}(16)$ | $3.12 \times 10^{8}$ | $6.07 \times 10^{8}$ | 10.14 | $1.03 \times 10^{9}$ | 5.99 | 1.37 |
| SBS1159+545 | $\leq 6.30 \times 10^{7 a}$ (5) | - | - | 7.44 (1) | $\leq 7.76 \times 10^{5 d}(-)$ | $\leq 4.26 \times 10^{7}$ | $\leq 8.74 \times 10^{7}$ | - | $\leq 1.45 \times 10^{8}$ | - | 1.37 |
| SBS1211+540 | $5.60 \times 10^{7 c}(12)$ | 16.08 | 0.1 | 7.58 (1) | $6.90 \times 10^{5 d}(-)$ | $3.79 \times 10^{7}$ | $7.78 \times 10^{7}$ | 15.89 | $1.29 \times 10^{8}$ | 9.59 | 1.37 |
| SBS1249+493 | $1.00 \times 10^{9 c}(25)$ | 48.15 | 0.06 | 7.68 (1) | $1.23 \times 10^{7 d}(-)$ | $6.77 \times 10^{8}$ | $1.39 \times 10^{9}$ | 47.56 | $2.30 \times 10^{9}$ | 28.71 | 1.37 |
| SBS1415+437 | $4.37 \times 10^{8 c}(12)$ | 16.08 | 0.4 | 7.55 (1) | $5.38 \times 10^{6 d}(-)$ | $2.96 \times 10^{8}$ | $6.06 \times 10^{8}$ | 15.89 | $1.00 \times 10^{9}$ | 9.59 | 1.37 |
| SBS1533+574 | $3.00 \times 10^{9 c}(14)$ | 11.29 | 0.2 | 8.05 (1) | $3.70 \times 10^{7 d}(-)$ | $2.03 \times 10^{9}$ | $4.18 \times 10^{9}$ | 11.15 | $6.92 \times 10^{9}$ | 6.73 | 1.38 |
| Tol0618-402 | - | - | - | 8.09 (1) | - | 8 |  | - | - | - | 1.38 |
| Tol1214-277 | $\leq 3.22 \times 10^{8 a}$ (5) | - | - | 7.52 (1) | $\leq 3.97 \times 10^{6 d}(-)$ | $\leq 2.18 \times 10^{8}$ | $\leq 4.47 \times 10^{8}$ | - | $\leq 7.41 \times 10^{8}$ | - | 1.37 |
| UGC4483 | $2.52 \times 10^{7 b}$ (26) | 5.00 | 14.4 | 7.46 (1) | $4.06 \times 10^{5 d}$ (14) | $2.23 \times 10^{7}$ | $3.51 \times 10^{7}$ | 4.92 | $6.52 \times 10^{7}$ | 2.65 | 1.37 |
| UGCA20 | $6.92 \times 10^{6 b}(27)$ | 13.82 | 9.8 | 7.50 (1) | $2.46 \times 10^{5 d}(-)$ | $1.35 \times 10^{7}$ | $9.83 \times 10^{6}$ | 13.34 | $2.80 \times 10^{7}$ | 4.68 | 1.37 |

Chapter 9. Gas-to-Dust mass ratios as a function of metallicity
Table 9.3. Gas masses for the DGS, KINGFISH and G11 samples (continued).

|  | $M_{H I}\left[M_{\odot}\right]$ |  |  | $M_{H_{2}}\left[M_{\odot}\right]$ |  |  | $\mathbf{M}_{\text {gas }}\left[M_{\odot}\right]$ |  |  |  |  |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Name | $M_{H I}$ (Ref) | Unc (\%) | HIarea ( $\mathrm{arcmin}^{2}$ ) | $12+\log (\mathrm{O} / \mathrm{H})$ (Ref) | $M_{H_{2}, M W}$ (Ref) | $M_{H_{2}, Z}$ | $\mathrm{M}_{\text {gas }, M W}$ | Unc (\%) | $\mathrm{M}_{\text {gas }, Z}$ | Unc (\%) | $\mu_{\text {gal }}$ |
| UM133 | $2.15 \times 10^{86}(28)$ | 4.38 | 6.2 | 7.82 (1) | $4.93 \times 10^{6 d}$ (-) | $2.71 \times 10^{8}$ | $3.02 \times 10^{8}$ | 4.28 | $6.67 \times 10^{8}$ | 1.94 | 1.37 |
| UM311 | $6.09 \times 10^{10 b}(29)$ | 16.08 | 21.2 | 8.36 (1) | $9.37 \times 10^{7}$ (1) | $4.28 \times 10^{8}$ | $8.44 \times 10^{10}$ | 16.06 | $8.48 \times 10^{10}$ | 15.97 | 1.38 |
| UM448 | $6.00 \times 10^{9 a}(15)$ | 27.63 | - | 8.32 (1) | $3.55 \times 10^{9}$ (3) | $1.92 \times 10^{10}$ | $1.32 \times 10^{10}$ | 17.36 | $3.48 \times 10^{10}$ | 6.59 | 1.38 |
| UM461 | $7.31 \times 10^{7 b}(26)$ | 5.00 | 2.6 | 7.73 (1) | $1.66 \times 10^{6 d}$ (3) | $9.09 \times 10^{7}$ | $1.03 \times 10^{8}$ | 4.89 | $2.25 \times 10^{8}$ | 2.23 | 1.37 |
| VIIZw403 | $3.25 \times 10^{7 b}(2)$ | 8.28 | 10.4 | 7.66 (1) | $6.41 \times 10^{5 d}(17)$ | $3.52 \times 10^{7}$ | $4.54 \times 10^{7}$ | 8.12 | $9.29 \times 10^{7}$ | 3.97 | 1.37 |
| KINGFISH |  |  |  |  |  |  |  |  |  |  |  |
| NGC0337 | $3.34 \times 10^{9}(27)$ | 11.97 | - | 8.18 (2) | $\leq 6.98 \times 10^{8}(18)$ | $\leq 7.31 \times 10^{9}$ | $\leq 5.57 \times 10^{9}$ | - | $\leq 1.47 \times 10^{10}$ | - | 1.38 |
| NGC0584 | $1.58 \times 10^{8}(27)$ | 35.92 | - | 8.43 (2) | $\leq 4.78 \times 10^{8}(18)$ | $\leq 1.58 \times 10^{9}$ | $\leq 8.82 \times 10^{8}$ | - | $\leq 2.41 \times 10^{9}$ | - | 1.39 |
| NGC0628 | $3.72 \times 10^{9}$ (30) | 16.08 | - | 8.35 (2) | $8.71 \times 10^{8}(19)$ | $4.17 \times 10^{9}$ | $6.35 \times 10^{9}$ | 13.04 | $1.09 \times 10^{10}$ | 7.59 | 1.38 |
| NGC0855 | $1.32 \times 10^{8}(27)$ | 29.47 | - | 8.29 (2) | $\leq 8.91 \times 10^{7}(18)$ | $\leq 5.62 \times 10^{8}$ | $\leq 3.05 \times 10^{8}$ | - | $\leq 9.59 \times 10^{8}$ | - | 1.38 |
| NGC0925 | $4.56 \times 10^{9}(30)$ | 16.08 | - | 8.25 (2) | $6.14 \times 10^{8}(19)$ | $4.66 \times 10^{9}$ | $7.15 \times 10^{9}$ | 14.18 | $1.27 \times 10^{10}$ | 7.96 | 1.38 |
| NGC1097 | $7.56 \times 10^{9}(30)$ | 16.08 | - | 8.47 (2) | $1.67 \times 10^{8}(20)$ | $4.59 \times 10^{8}$ | $1.07 \times 10^{10}$ | 15.74 | $1.11 \times 10^{10}$ | 15.16 | 1.39 |
| NGC1266 | .56×10 (30) | - | - | 8.29 (2) | $1.33 \times 10^{9}(21)$ | $8.39 \times 10^{9}$ | $\geq 1.84 \times 10^{9}$ | - | $\geq 1.16 \times 10^{10}$ | - | 1.38 |
| NGC1291 | $1.78 \times 10^{9}(30)$ | 16.08 | - | 8.52 (2) | - | - | $\geq 2.47 \times 10^{9}$ | - | $\geq 2.47 \times 10^{9}$ | - | 1.39 |
| NGC1316 | $\leq 4.73 \times 10^{8}$ (30) | - | - | 8.77 (2) | - | - | - | - | - | - | 1.40 |
| NGC1377 | - | - | - | 8.29 (2) | - | - | - | - | - | - | 1.38 |
| NGC1404 | - | - | - | 8.54 (2) | - | - | - | - | - | - | 1.39 |
| IC0342 | $9.49 \times 10^{9}$ (31) | 16.08 | - | 8.46 (3) | $1.60 \times 10^{9}(20)$ | $4.60 \times 10^{9}$ | $1.54 \times 10^{10}$ | 13.77 | $1.95 \times 10^{10}$ | 10.83 | 1.39 |
| NGC1482 | $\leq 8.01 \times 10^{8}$ (30) | - | - | 8.11 (2) | $3.11 \times 10^{9}$ (19) | $4.49 \times 10^{10}$ | $\leq 5.38 \times 10^{9}$ | - | $\leq 6.30 \times 10^{10}$ | - | 1.38 |
| NGC1512 | $7.33 \times 10^{9}(30)$ | 16.08 | - | 8.56 (2) | ${ }^{-}$ | - | $\geq 1.02 \times 10^{10}$ | - | $\geq 1.02 \times 10^{10}$ | - | 1.39 |
| NGC2146 | $3.89 \times 10^{9}(27)$ | 22.10 | - | 8.68 (2) | $7.77 \times 10^{10}(22)$ | $8.13 \times 10^{10}$ | $1.14 \times 10^{11}$ | 1.05 | $1.19 \times 10^{11}$ | 1.01 | 1.40 |
| HoII | $4.21 \times 10^{8}(27)$ | 23.95 | - | 7.72 (2) | $5.18 \times 10^{6 d}(18)$ | $2.85 \times 10^{8}$ | $5.85 \times 10^{8}$ | 23.66 | $9.69 \times 10^{8}$ | 14.28 | 1.37 |
| DDO053 | $6.16 \times 10^{7}(30)$ | 16.08 | - | 7.60 (2) | $7.60 \times 10^{5 d}$ (11) | $4.17 \times 10^{7}$ | $8.56 \times 10^{7}$ | 15.89 | $1.42 \times 10^{8}$ | 9.59 | 1.37 |
| NGC2798 | $2.13 \times 10^{9}$ (30) | 16.08 | - | 8.34 (2) | $3.21 \times 10^{9}$ (19) | $1.61 \times 10^{10}$ | $7.38 \times 10^{9}$ | 6.41 | $2.52 \times 10^{10}$ | 1.88 | 1.38 |
| NGC2841 | $8.63 \times 10^{9}(30)$ | 16.08 | - | 8.54 (2) | $2.92 \times 10^{9}$ (19) | $5.82 \times 10^{9}$ | $1.60 \times 10^{10}$ | 12.02 | $2.01 \times 10^{10}$ | 9.61 | 1.39 |
| NGC2915 | $3.52 \times 10^{8}(30)$ | 16.08 | - | 7.94 (2) | $4.34 \times 10^{6 d}(-)$ | $2.39 \times 10^{8}$ | $4.90 \times 10^{8}$ | 15.89 | $8.12 \times 10^{8}$ | 9.59 | 1.38 |
| HoI | $1.45 \times 10^{8}(16)$ | 16.08 | - | 7.61 (2) | $1.79 \times 10^{6 d}$ (18) | $9.85 \times 10^{7}$ | $2.02 \times 10^{8}$ | 15.89 | $3.35 \times 10^{8}$ | 9.59 | 1.37 |
| NGC2976 | $1.27 \times 10^{8}(30)$ | 16.08 | - | 8.36 (2) | $5.81 \times 10^{7}(19)$ | $2.66 \times 10^{8}$ | $2.57 \times 10^{8}$ | 11.05 | $5.44 \times 10^{8}$ | 5.21 | 1.38 |
| NGC3049 | $1.21 \times 10^{9}(30)$ | 16.08 | - | 8.53 (2) | $1.80 \times 10^{8}(18)$ | $3.75 \times 10^{8}$ | $1.93 \times 10^{9}$ | 14.00 | $2.20 \times 10^{9}$ | 12.27 | 1.39 |
| NGC3077 | $8.71 \times 10^{8}(16)$ | 16.08 | - | 8.52 (4) | $3.80 \times 10^{6}$ (9) | $8.32 \times 10^{6}$ | $1.21 \times 10^{9}$ | 16.01 | $1.22 \times 10^{9}$ | 15.93 | 1.39 |
| M81dwB | $1.16 \times 10^{7}(16)$ | 16.08 | - | 7.84 (2) | $1.43 \times 10^{5 d}$ (18) | $7.85 \times 10^{6}$ | $1.61 \times 10^{7}$ | 15.89 | $2.67 \times 10^{7}$ | 9.59 | 1.37 |
| NGC3190 | $4.27 \times 10^{8}(27)$ | 36.84 | - | 8.49 (2) | $\leq 3.93 \times 10^{8}(18)$ | $\leq 9.87 \times 10^{8}$ | $\leq 1.14 \times 10^{9}$ | - | $\leq 1.96 \times 10^{9}$ | - | 1.39 |
| NGC3184 | $3.37 \times 10^{9}(30)$ | 16.08 | - | 8.51 (2) | $1.19 \times 10^{9}$ (19) | $2.73 \times 10^{9}$ | $6.33 \times 10^{9}$ | 11.88 | $8.47 \times 10^{9}$ | 8.88 | 1.39 |
| NGC3198 | $6.85 \times 10^{9}(27)$ | 26.71 | - | 8.34 (2) | $1.30 \times 10^{9}(18)$ | $6.53 \times 10^{9}$ | $1.13 \times 10^{10}$ | 22.44 | $1.85 \times 10^{10}$ | 13.68 | 1.38 |
| IC2574 | $1.32 \times 10^{9}$ (30) | 16.08 | - | 7.85 (2) | $3.16 \times 10^{6}$ (19) | $1.51 \times 10^{8}$ | $1.82 \times 10^{9}$ | 16.05 | $2.02 \times 10^{9}$ | 14.43 | 1.37 |
| NGC3265 | $1.75 \times 10^{8}(27)$ | 38.68 | - | 8.27 (2) | $\leq 4.39 \times 10^{8}(18)$ | $\leq 3.04 \times 10^{9}$ | $\leq 8.47 \times 10^{8}$ | - | $\leq 4.43 \times 10^{9}$ | - | 1.38 |
| NGC3351 | $1.02 \times 10^{9}(30)$ | 16.08 | - | 8.60 (2) | $4.77 \times 10^{8}(19)$ | $7.23 \times 10^{8}$ | $2.09 \times 10^{9}$ | 10.97 | $2.43 \times 10^{9}$ | 9.43 | 1.39 |
| NGC3521 | $8.71 \times 10^{9}(30)$ | 16.08 | - | 8.39 (2) | $4.77 \times 10^{9}(19)$ | $1.90 \times 10^{10}$ | $1.87 \times 10^{10}$ | 10.39 | $3.84 \times 10^{10}$ | 5.05 | 1.38 |
| NGC3621 | $6.99 \times 10^{9}(30)$ | 16.08 | - | 8.27 (2) | ( | - | $\geq 9.65 \times 10^{9}$ | - | $\geq 9.65 \times 10^{9}$ | - | 1.38 |
| NGC3627 | $8.46 \times 10^{8}(30)$ | 16.08 | - | 8.34 (2) | $3.21 \times 10^{9}(19)$ | $1.61 \times 10^{10}$ | $5.61 \times 10^{9}$ | 3.36 | $2.34 \times 10^{10}$ | 0.80 | 1.38 |
| NGC3773 | $9.03 \times 10^{7}(30)$ | 16.08 | - | 8.43 (2) | $4.95 \times 10^{7}(18)$ | $1.64 \times 10^{8}$ | $1.94 \times 10^{8}$ | 10.39 | $3.52 \times 10^{8}$ | 5.71 | 1.39 |
| NGC3938 | $8.00 \times 10^{9}(30)$ | 16.08 | - | 8.42 (2) | $4.38 \times 10^{9}$ (19) | $1.52 \times 10^{10}$ | $1.71 \times 10^{10}$ | 10.39 | $3.21 \times 10^{10}$ | 5.54 | 1.38 |
| NGC4236 | $2.69 \times 10^{9}(30)$ | 16.08 | - | 8.17 (2) | $9.33 \times 10^{7}(18)$ | $1.02 \times 10^{9}$ | $3.84 \times 10^{9}$ | 15.55 | $5.12 \times 10^{9}$ | 11.65 | 1.38 |

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|  | $M_{H I}\left[M_{\odot}\right]$ |  |  | $M_{H_{2}}\left[M_{\odot}\right]$ |  |  | $\mathbf{M}_{\text {gas }}$ [ $M_{\odot}$ ] |  |  |  |  |
| :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: | :---: |
| Name | $M_{H I}$ (Ref) | Unc (\%) | HIarea ( $\mathrm{arcmin}^{2}$ ) | $12+\log (\mathrm{O} / \mathrm{H})($ Ref $)$ | $M_{H_{2}, M W}$ (Ref) | $M_{H_{2}, Z}$ | $\mathrm{M}_{\text {gas }, \text { M }}$ | Unc (\%) | $\mathrm{M}_{\text {gas }, Z}$ | Unc (\%) | $\mu_{\text {gal }}$ |
| NGC4254 | $3.76 \times 10^{9}{ }^{(30)}$ | 16.08 | - | 8.45 (2) | $6.82 \times 10^{9}$ (19) | $2.06 \times 10^{10}$ | $1.47 \times 10^{10}$ | 5.71 | $3.37 \times 10^{10}$ | 2.48 | 1.39 |
| NGC4321 | $2.39 \times 10^{9}(30)$ | 16.08 | - | 8.50 (2) | $5.34 \times 10^{9}(19)$ | $1.28 \times 10^{10}$ | $1.07 \times 10^{10}$ | 4.97 | $2.11 \times 10^{10}$ | 2.53 | 1.39 |
| NGC4536 | $1.73 \times 10^{9}(30)$ | 16.08 | - | 8.21 (2) | $1.72 \times 10^{9}$ (19) | $1.57 \times 10^{10}$ | $4.75 \times 10^{9}$ | 8.05 | $2.40 \times 10^{10}$ | 1.59 | 1.38 |
| NGC4559 | $4.09 \times 10^{9}(30)$ | 16.08 | - | 8.29 (2) | $1.92 \times 10^{8}$ (18) | $1.21 \times 10^{9}$ | $5.91 \times 10^{9}$ | 15.36 | $7.32 \times 10^{9}$ | 12.41 | 1.38 |
| NGC4569 | $1.55 \times 10^{8}(30)$ | 16.08 | - | 8.58 (2) | $1.14 \times 10^{9}$ (19) | $1.90 \times 10^{9}$ | $1.81 \times 10^{9}$ | 1.92 | $2.86 \times 10^{9}$ | 1.21 | 1.39 |
| NGC4579 | $5.47 \times 10^{8}(30)$ | 16.08 | - | 8.54 (2) | $1.89 \times 10^{9}(19)$ | $3.77 \times 10^{9}$ | $3.39 \times 10^{9}$ | 3.61 | $6.00 \times 10^{9}$ | 2.04 | 1.39 |
| NGC4594 | $2.60 \times 10^{8}(30)$ | 16.08 | - | 8.54 (2) | $2.16 \times 10^{8}$ (19) | $4.31 \times 10^{8}$ | $6.62 \times 10^{8}$ | 8.79 | $9.61 \times 10^{8}$ | 6.06 | 1.39 |
| NGC4625 | $1.00 \times 10^{9}(30)$ | 16.08 | - | 8.35 (2) | $2.09 \times 10^{7}(18)$ | $1.00 \times 10^{8}$ | $1.42 \times 10^{9}$ | 15.76 | $1.53 \times 10^{9}$ | 14.62 | 1.38 |
| NGC4631 | $8.77 \times 10^{9}(30)$ | 16.08 | - | 8.12 (2) | $1.10 \times 10^{9}(19)$ | $1.52 \times 10^{10}$ | $1.36 \times 10^{10}$ | 14.29 | $3.30 \times 10^{10}$ | 5.88 | 1.38 |
| NGC4725 | $3.59 \times 10^{9}(30)$ | 16.08 | - | 8.35 (2) | $2.16 \times 10^{9}(19)$ | $1.03 \times 10^{10}$ | $7.95 \times 10^{9}$ | 10.05 | $1.92 \times 10^{10}$ | 4.15 | 1.38 |
| NGC4736 | $4.03 \times 10^{8}(30)$ | 16.08 | - | 8.31 (2) | $4.41 \times 10^{8}(19)$ | $2.54 \times 10^{9}$ | $1.17 \times 10^{9}$ | 7.68 | $4.06 \times 10^{9}$ | 2.20 | 1.38 |
| DDO154 | $3.55 \times 10^{8}(16)$ | 16.08 | - | 7.54 (2) | $4.37 \times 10^{6 d}(23)$ | $2.40 \times 10^{8}$ | $4.93 \times 10^{8}$ | 15.89 | $8.17 \times 10^{8}$ | 9.59 | 1.37 |
| NGC4826 | $2.77 \times 10^{8}(30)$ | 16.08 | - | 8.54 (2) | $4.80 \times 10^{8}(19)$ | $9.58 \times 10^{8}$ | $1.05 \times 10^{9}$ | 5.88 | $1.72 \times 10^{9}$ | 3.61 | 1.39 |
| DDO165 | $1.12 \times 10^{8}(27)$ | 14.74 | - | 7.63 (2) | $1.38 \times 10^{6 d}$ (18) | $7.60 \times 10^{7}$ | $1.56 \times 10^{8}$ | 14.56 | $2.58 \times 10^{8}$ | 8.79 | 1.37 |
| NGC5055 | $5.59 \times 10^{9}{ }^{(30)}$ | 16.08 | - | 8.40 (2) | $2.73 \times 10^{9}$ (19) | $1.04 \times 10^{10}$ | $1.15 \times 10^{10}$ | 10.80 | $2.21 \times 10^{10}$ | 5.63 | 1.38 |
| NGC5398 | $2.48 \times 10^{8}(27)$ | 11.97 | - | 8.35 (2) | ( | - | $\geq 3.43 \times 10^{8}$ | - | $\geq 3.43 \times 10^{8}$ | - | 1.38 |
| NGC5408 | $3.22 \times 10^{8}(30)$ | 16.08 | - | 7.81 (2) | $3.97 \times 10^{6 d}(-)$ | $2.18 \times 10^{8}$ | $4.48 \times 10^{8}$ | 15.89 | $7.43 \times 10^{8}$ | 9.59 | 1.37 |
| NGC5457 | $1.16 \times 10^{10}(16)$ | 16.08 | - | 8.42 (5) | -97× | - | $\geq 1.60 \times 10^{10}$ | - | $\geq 1.60 \times 10^{10}$ | - | 1.38 |
| NGC5474 | $9.71 \times 10^{8}(27)$ | 24.87 | - | 8.31 (2) | $\leq 5.88 \times 10^{7}$ (18) | $\leq 3.38 \times 10^{8}$ | $\leq 1.42 \times 10^{9}$ | - | $\leq 1.81 \times 10^{9}$ | - | 1.38 |
| NGC5713 | $5.51 \times 10^{9}(30)$ | 16.08 | - | 8.24 (2) | ${ }^{5} .39 \times 10^{9}$ (19) | $2.69 \times 10^{10}$ | $1.23 \times 10^{10}$ | 9.96 | $4.47 \times 10^{10}$ | 2.73 | 1.38 |
| NGC5866 | $\leq 2.85 \times 10^{8}$ (30) | - | - | 8.47 (2) | $6.38 \times 10^{8}(19)$ | $1.76 \times 10^{9}$ | $\leq 1.28 \times 10^{9}$ | - | $\leq 2.83 \times 10^{9}$ | - | 1.39 |
| NGC6946 | $3.58 \times 10^{9}(30)$ | 16.08 | - | 8.40 (2) | $4.50 \times 10^{9}(19)$ | $1.71 \times 10^{10}$ | $1.12 \times 10^{10}$ | 7.13 | $2.86 \times 10^{10}$ | 2.79 | 1.38 |
| NGC7331 | $8.85 \times 10^{9}(30)$ | 16.08 | - | 8.34 (2) | $6.70 \times 10^{9}(19)$ | $3.36 \times 10^{10}$ | $2.15 \times 10^{10}$ | 9.16 | $5.86 \times 10^{10}$ | 3.36 | 1.38 |
| NGC7793 | $8.75 \times 10^{8}(30)$ | 16.08 | - | 8.31 (2) | - | - | $\geq 1.21 \times 10^{9}$ | - | $\geq 1.21 \times 10^{9}$ | - | 1.38 |
| G11 |  |  |  |  |  |  |  |  |  |  |  |
| M83 | $5.13 \times 10^{9}(31)$ | 5.00 | - | 8.62 (6) | $8.65 \times 10^{8}(20)$ | $1.19 \times 10^{9}$ | $8.35 \times 10^{9}$ | 4.28 | $8.81 \times 10^{9}$ | 4.06 | 1.39 |
| NGC1808 | $1.78 \times 10^{9}(31)$ | 9.46 | - | 9.10 (7) | $1.75 \times 10^{9}(20)$ | $2.64 \times 10^{8}$ | $5.08 \times 10^{9}$ | 4.77 | $2.95 \times 10^{9}$ | 8.24 | 1.44 |
| NGC7552 | $4.79 \times 10^{9}(31)$ | 16.08 | - | 8.35 (8) | 175x10 (20) | - | $\geq 6.62 \times 10^{9}$ | - | $\geq 6.62 \times 10^{9}$ | - | 1.38 |
| M82 | $8.91 \times 10^{8}(31)$ | 10.29 | - | 8.51 (8) | $4.43 \times 10^{8}(20)$ | $1.01 \times 10^{9}$ | $1.85 \times 10^{9}$ | 6.87 | $2.65 \times 10^{9}$ | 4.81 | 1.39 |
| NGC1068 | $2.24 \times 10^{9}(31)$ | 18.35 | - | 9.00 (7) | $7.26 \times 10^{9}(20)$ | $1.74 \times 10^{9}$ | $1.36 \times 10^{10}$ | 4.32 | $5.68 \times 10^{9}$ | 10.32 | 1.43 |
| NGC0891 | $7.59 \times 10^{9}(31)$ | 16.08 | - | 8.90 (7) | $5.75 \times 10^{9}(20)$ | $2.19 \times 10^{9}$ | $1.89 \times 10^{10}$ | 9.15 | $1.38 \times 10^{10}$ | 12.48 | 1.42 |
| MGC+02-04-025 | $1.74 \times 10^{9}(31)$ | 16.08 | - | 8.52 (9) | - | - | $\geq 2.41 \times 10^{9}$ | - | $\geq 2.41 \times 10^{9}$ | - | 1.39 |
| NGC7469 | $1.51 \times 10^{9}(31)$ | 16.08 | - | 8.70 (10) | $6.28 \times 10^{9}(20)$ | $6.00 \times 10^{9}$ | $1.09 \times 10^{10}$ | 3.12 | $1.05 \times 10^{10}$ | 3.24 | 1.40 |
| NGC5256 | - | - | - | 8.53 (9) | $9.92 \times 10^{9}$ (20) | $2.07 \times 10^{10}$ | $\geq 1.38 \times 10^{10}$ | - | $\geq 2.88 \times 10^{10}$ | - | 1.39 |
| NGC5953 | $5.75 \times 10^{8}(31)$ | 16.08 | - | 8.73 (7) | $1.62 \times 10^{9}(20)$ | $1.34 \times 10^{9}$ | $3.07 \times 10^{9}$ | 4.22 | $2.69 \times 10^{9}$ | 4.82 | 1.40 |
| M51 | $5.01 \times 10^{9}(31)$ | 5.63 | - | 8.55 (8) | $3.95 \times 10^{9}(20)$ | $7.53 \times 10^{9}$ | $1.25 \times 10^{10}$ | 3.15 | $1.74 \times 10^{10}$ | 2.25 | 1.39 |
| NGC3995 | $6.17 \times 10^{9}(31)$ | 16.08 | - | 8.66 (7) | - | - | $\geq 8.61 \times 10^{9}$ | - | $\geq 8.61 \times 10^{9}$ | - | 1.40 |
| NGC3994 | $2.82 \times 10^{9}(31)$ | 16.08 | - | 8.61 (7) | ${ }^{-}$ | - | $\geq 3.93 \times 10^{9}$ | - | $\geq 3.93 \times 10^{9}$ | - | 1.39 |
| NGC6052 | $3.80 \times 10^{9}(31)$ | 16.08 | - | 8.65 (7) | $3.00 \times 10^{9}(20)$ | $3.61 \times 10^{9}$ | $9.49 \times 10^{9}$ | 8.99 | $1.03 \times 10^{10}$ | 8.25 | 1.40 |
| NGC1222 | $1.20 \times 10^{9}(31)$ | 16.08 | - | 8.57 (7) | $7.23 \times 10^{8}(21)$ | $1.26 \times 10^{9}$ | $2.68 \times 10^{9}$ | 10.04 | $3.42 \times 10^{9}$ | 7.86 | 1.39 |
| NGC7674 | $1.07 \times 10^{10}(31)$ | 16.08 | - | 8.14 (9) | $6.34 \times 10^{9}(20)$ | $7.99 \times 10^{10}$ | $2.35 \times 10^{10}$ | 10.10 | $1.25 \times 10^{11}$ | 1.90 | 1.38 |
| NGC4670 | $1.66 \times 10^{8}(31)$ | 16.08 | - | 8.30 (7) | $9.00 \times 10^{6}(20)$ | $5.42 \times 10^{7}$ | $2.42 \times 10^{8}$ | 15.26 | $3.04 \times 10^{8}$ | 12.12 | 1.38 |

References for metallicities (1) Madden et al. (2013) ; (2) Kennicutt et al. (2011) ; (3) McCall et al. (1985) ; (4) McQuade et al. (1995) ; (5) Kennicutt et al. (2003b) (6) Pilyugin et al. (2006) ; (7) Galametz et al. (2011) ; (8) Moustakas et al. (2010) ; (9) Veilleux et al. (1995) ; (10) Bonatto \& Pastoriza (1990)

References for HI masses : (1) Cormier et al. (2014) ; (2) Thuan et al. (2004) ; (3) Gordon \& Gottesman (1981) ; (4) Sauvage et al. (1997) ; (5) Pustilnik \& Martin (2007) ; (6) MacHattie et al., in prep. ; (7) Chengalur et al. (2006) ; (8) Lelli et al. (2012) ; (9) Huchtmeier \& Richter (1988) ; (10) Bettoni et al. (2003) ; (11) Williams et al. (1991) ; (12) Huchtmeier et al. (2005) ; (13) Thuan \& Martin (1981) ; (14) Thuan et al. (1999a) ; (15) Davoust \& Contini (2004) ; (16) Walter et al. (2008) ; (17) Meurer et al. (1998) ; (18) Hunter et al. (2011) ; (19) van Eymeren et al. (2009) ; (20) López-Sánchez et al. (2012) ; (21) Cannon et al. (2004) ; (22) de Blok \& Walter (2006) ; (23) Begum \& Chengalur (2005b) ; (24) Ekta et al. (2009) ; (25) Pustilnik et al. (2002) ; (26) van Zee et al. (1998) ; (27) Paturel et al. (2003) ; (28) Ekta \& Chengalur (2010) ; (29) Moles et al. (1994) ; (30) Draine et al. (2007) ; (31) Galametz et al. (2011)

References for $\mathbf{H}_{2}$ masses (1) Cormier et al. (2014) ; (2) Israel (2005) ; (3) Sage et al. (1992) ; (4) Kobulnicky et al. (1995) ; (5) Leroy et al. (2007) ; (6) Leroy et al. (2006) ; (7) Bettoni et al. (2003) ; (8) Young et al. (1995) ; (9) Albrecht et al. (2004) ; (10) Greve et al. (1996) ; (11) Schruba et al. (2012) ; (12) Walter et al. (2001) ; (13) Böttner et al. (2003) ; (14) Taylor et al. (1998) ; (15) Gratier et al. (2010) ; (16) Dale et al. (2001b) ; (17) Leroy et al. (2005) ; (18) Wilson et al. (2012) ; (19) Draine et al. (2007) ; (20) Galametz et al. (2011) ; (21) Young et al. (2011) ; (22) Young \& Knezek (1989) ; (23) Leroy et al. (2009)

Note on mass references : The references given here are where the original measurement was found. If a correction was applied, the quoted masses are the corrected masses (see ${ }^{b},{ }^{c},{ }^{d}$ ) and not the original masses. If no correction was applied, the quoted masses are the original masses. Note that some galaxies have a molecular gas mass but no reference : this means that the molecular gas mass has been corrected (see ${ }^{d}$ ) but that no measurements of the molecular gas mass is available in the literature.
${ }^{a}$ For these galaxies, no information was available on the extent of the HI observations. The quoted HI mass is thus the total HI mass (diamonds in Fig. 9.4). ${ }^{b}$ For these galaxies, we were able to correct the Hı mass and to rescale it to the Hı mass contained in the IR aperture (triangles in Fig. 9.4).
${ }^{c}$ For these galaxies, the Hi extent is similar to the IR aperture size and we consider that the total Hi mass corresponds to the Hi mass within the IR aperture (crosses in Fig. 9.4).
${ }^{c 2}$ For these galaxies, the H mass corresponding to the dust emitting region is given in the reference, and is the mass quoted here. ${ }^{d}$ For these galaxies, the $\mathrm{H}_{2}$ mass correction was applied : i.e. non-detections or absence of data were replaced by $0.012 \times M_{H I}$ for for galaxies with $12+\log (\mathrm{O} / \mathrm{H}) \leq 8.1$. ${ }^{e}$ For these 2 galaxies, the Hir mass has a significant contribution to the total gas mass : $M_{\text {HII }} \sim 1.2 \times M_{H I}$ for Haro 11 from for Pox 186 from Gil de Paz et al. (2003) in the $X_{C O, M W}$ case, and has to be taken into account to estimate the total gas mass.

Table 9.3. G/D for the DGS, KINGFISH and G11 samples.

| Name | $\mathrm{G} / \mathrm{D}\left(X_{C O, M W}\right)$ | Uncertainty (\%) | $\mathrm{G} / \mathrm{D}\left(X_{C O, Z}\right)$ | Uncertainty (\%) | Comments ${ }^{\text {a }}$ |
| :---: | :---: | :---: | :---: | :---: | :---: |
| DGS |  |  |  |  |  |
| Haro11 | $1.75 \times 10^{2}$ | 21.77 | $3.04 \times 10^{2}$ | 14.43 | - |
| Haro2 | $3.97 \times 10^{2}$ | 16.43 | $1.11 \times 10^{3}$ | 11.96 | - |
| Haro3 | $7.08 \times 10^{2}$ | 12.27 | $8.12 \times 10^{2}$ | 12.22 | - |
| He2-10 | $3.54 \times 10^{2}$ | 12.07 | $5.51 \times 10^{2}$ | 11.46 | - |
| HS0017+1055 | $\leq 2.93 \times 10^{3}$ | - | $\leq 4.86 \times 10^{3}$ | - | $1+2$ |
| HS0052+2536 | $\leq 1.02 \times 10^{4}$ | - | $\leq 1.70 \times 10^{4}$ | - | $1+2$ |
| HS0822+3542 | $1.29 \times 10^{4}$ | 54.02 | $2.42 \times 10^{4}$ | 50.76 | 3 |
| HS1222+3741 | - | - | - | - | - |
| HS1236+3937 | - | - | - | - | - |
| HS1304+3529 | - | - | - | - | - |
| HS1319+3224 | - | - | - | - | - |
| HS1330+3651 | - | - | - | - | - |
| HS1442+4250 | - | - | - | - | - |
| HS2352+2733 | - | - | - | - | - |
| IZw18 | $2.76 \times 10^{5}$ | 23.57 | $6.69 \times 10^{5}$ | 18.77 | 3 |
| IC10 | $2.40 \times 10^{2}$ | 34.26 | $4.76 \times 10^{2}$ | 34.08 | - |
| IIZw40 | $1.23 \times 10^{3}$ | 46.42 | $1.67 \times 10^{3}$ | 45.25 | - |
| Mrk1089 | $7.93 \times 10^{2}$ | 15.47 | $1.01 \times 10^{3}$ | 14.72 | - |
| Mrk1450 | $1.19 \times 10^{3}$ | 208.33 | $1.97 \times 10^{3}$ | 207.94 | - |
| Mrk153 | $\leq 5.62 \times 10^{3}$ | - | $\leq 9.25 \times 10^{3}$ | - | $1+2$ |
| Mrk209 | $9.15 \times 10^{3}$ | 39.03 | $8.26 \times 10^{4}$ | 38.58 | - |
| Mrk930 | $7.38 \times 10^{2}$ | 55.41 | $8.82 \times 10^{2}$ | 55.08 | - |
| NGC1140 | $1.23 \times 10^{3}$ | 29.20 | $1.26 \times 10^{3}$ | 28.79 | - |
| NGC1569 | $8.55 \times 10^{2}$ | 22.39 | $9.25 \times 10^{2}$ | 21.53 | - |
| NGC1705 | $\geq 8.24 \times 10^{2}$ | - | $\geq 8.24 \times 10^{2}$ | - | 2 |
| NGC2366 | $3.60 \times 10^{3}$ | 34.47 | $9.47 \times 10^{3}$ | 33.93 | - |
| NGC4214 | $3.14 \times 10^{2}$ | 30.52 | $3.22 \times 10^{2}$ | 30.30 | - |
| NGC4449 | $4.46 \times 10^{2}$ | 18.79 | $5.48 \times 10^{2}$ | 16.45 | - |
| NGC4861 | $9.52 \times 10^{2}$ | 26.80 | $1.68 \times 10^{3}$ | 26.31 | - |
| NGC5253 | $2.71 \times 10^{2}$ | 23.11 | $3.19 \times 10^{2}$ | 22.93 | - |
| NGC625 | $3.49 \times 10^{2}$ | 24.13 | $4.49 \times 10^{2}$ | 21.47 | - |
| NGC6822 | $6.13 \times 10^{2}$ | 112.72 | $6.73 \times 10^{2}$ | 112.64 | - |
| Pox186 | $\leq 2.49 \times 10^{1}$ | - | $\leq 3.45 \times 10^{1}$ | - | $1+2+3$ |
| SBS0335-052 | $1.95 \times 10^{5}$ | 143.38 | $3.30 \times 10^{5}$ | 143.15 | - |
| SBS1159+545 | $\leq 2.91 \times 10^{3}$ | - | $\leq 4.83 \times 10^{3}$ | - | $1+2+3$ |
| SBS1211+540 | $4.22 \times 10^{3}$ | 50.54 | $6.99 \times 10^{3}$ | 48.93 | - |
| SBS1249+493 | $3.31 \times 10^{3}$ | 57.66 | $5.49 \times 10^{3}$ | 43.44 | 3 |
| SBS1415+437 | $9.52 \times 10^{4}$ | 24.11 | $1.58 \times 10^{5}$ | 20.51 | 3 |
| SBS1533+574 | $2.11 \times 10^{3}$ | 24.14 | $3.49 \times 10^{3}$ | 22.44 | - |
| Tol0618-402 | - | - | - | - | - |
| Tol1214-277 | $\leq 3.57 \times 10^{3}$ | - | $\leq 5.92 \times 10^{3}$ | - | $1+2+3$ |
| UGC4483 | $9.45 \times 10^{4}$ | 41.33 | $1.75 \times 10^{5}$ | 41.13 | 3 |
| UGCA20 | - | - | - | - | 3 |
| UM133 | $3.16 \times 10^{3}$ | 58.97 | $6.99 \times 10^{3}$ | 58.85 | - |
| UM311 | $4.28 \times 10^{3}$ | 24.78 | $4.31 \times 10^{3}$ | 24.72 | - |
| UM448 | $7.36 \times 10^{2}$ | 19.27 | $1.94 \times 10^{3}$ | 10.64 | - |
| UM461 | $4.88 \times 10^{3}$ | 48.49 | $1.07 \times 10^{4}$ | 48.29 | - |
| VIIZw403 | $6.29 \times 10^{3}$ | 18.09 | $1.29 \times 10^{4}$ | 16.64 | - |
| KINGFISH |  |  |  |  |  |
| NGC0337 | $\leq 3.48 \times 10^{2}$ | - | $\leq 9.18 \times 10^{2}$ | - | 2 |
| NGC0584 | - | - | - | - | $2+3$ |
| NGC0628 | $2.47 \times 10^{2}$ | 18.01 | $4.24 \times 10^{2}$ | 14.56 | - |
| NGC0855 | $\leq 1.66 \times 10^{2}$ | - | $\leq 5.20 \times 10^{2}$ | - | 2 |
| NGC0925 | $2.84 \times 10^{2}$ | 20.23 | $5.06 \times 10^{2}$ | 16.49 | - |
| NGC1097 | $1.29 \times 10^{2}$ | 17.68 | $1.34 \times 10^{2}$ | 17.17 | - |
| NGC1266 | $\geq 2.18 \times 10^{2}$ | - | $\geq 1.37 \times 10^{3}$ | - | - |
| NGC1291 | $\geq 1.51 \times 10^{2}$ | - | $\geq 1.51 \times 10^{2}$ | - | 2 |

Table 9.3. G/D for the DGS, KINGFISH and G11 samples (continued).

| Name | $\mathrm{G} / \mathrm{D}\left(X_{C O, M W}\right)$ | Uncertainty (\%) | $\mathrm{G} / \mathrm{D}\left(X_{C O, Z}\right)$ | Uncertainty (\%) | Comments ${ }^{a}$ |
| :---: | :---: | :---: | :---: | :---: | :---: |
| NGC1316 | - | - | - | - | - |
| NGC1377 | - | - | - | - | - |
| NGC1404 | - | - | - | - | - |
| IC0342 | $3.89 \times 10^{2}$ | 15.60 | $4.94 \times 10^{2}$ | 13.08 | - |
| NGC1482 | $\leq 2.12 \times 10^{2}$ | - | $\leq 2.48 \times 10^{3}$ | - | 1 |
| NGC1512 | $\geq 3.29 \times 10^{2}$ | - | $\geq 3.29 \times 10^{2}$ | - | 2 |
| NGC2146 | $2.40 \times 10^{3}$ | 8.30 | $2.51 \times 10^{3}$ | 8.29 | - |
| HoII | $7.72 \times 10^{3}$ | 24.94 | $1.28 \times 10^{4}$ | 16.32 | - |
| DDO053 | $8.25 \times 10^{3}$ | 20.77 | $1.37 \times 10^{4}$ | 16.46 | - |
| NGC2798 | $5.94 \times 10^{2}$ | 9.39 | $2.03 \times 10^{3}$ | 7.11 | - |
| NGC2841 | $2.71 \times 10^{2}$ | 15.08 | $3.40 \times 10^{2}$ | 13.24 | - |
| NGC2915 | $1.45 \times 10^{3}$ | 26.39 | $2.40 \times 10^{3}$ | 23.15 | - |
| HoI | $8.50 \times 10^{2}$ | 17.91 | $1.41 \times 10^{3}$ | 12.66 | - |
| NGC2976 | $1.35 \times 10^{2}$ | 14.29 | $2.87 \times 10^{2}$ | 10.46 | - |
| NGC3049 | $2.51 \times 10^{2}$ | 24.91 | $2.86 \times 10^{2}$ | 23.98 | - |
| NGC3077 | $1.07 \times 10^{3}$ | 22.18 | $1.08 \times 10^{3}$ | 22.12 | - |
| M81dwB | $5.75 \times 10^{2}$ | 65.93 | $9.52 \times 10^{2}$ | 64.70 | - |
| NGC3190 | $\leq 6.26 \times 10^{1}$ | - | $\leq 1.08 \times 10^{2}$ | - | 2 |
| NGC3184 | $2.00 \times 10^{2}$ | 17.48 | $2.68 \times 10^{2}$ | 15.60 | - |
| NGC3198 | $3.00 \times 10^{2}$ | 26.10 | $4.92 \times 10^{2}$ | 19.09 | - |
| IC2574 | $6.53 \times 10^{2}$ | 50.08 | $7.26 \times 10^{2}$ | 49.59 | - |
| NGC3265 | $\leq 4.60 \times 10^{2}$ | - | $\leq 2.41 \times 10^{3}$ | - | 2 |
| NGC3351 | $1.08 \times 10^{2}$ | 12.91 | $1.26 \times 10^{2}$ | 11.63 | - |
| NGC3521 | $2.63 \times 10^{2}$ | 15.93 | $5.42 \times 10^{2}$ | 13.09 | - |
| NGC3621 | $\geq 5.10 \times 10^{2}$ | - | $\geq 5.10 \times 10^{2}$ | - | 2 |
| NGC3627 | $1.39 \times 10^{2}$ | 8.24 | $5.80 \times 10^{2}$ | 7.57 | - |
| NGC3773 | $2.56 \times 10^{2}$ | 16.22 | $4.66 \times 10^{2}$ | 13.71 | - |
| NGC3938 | $3.69 \times 10^{2}$ | 16.81 | $6.91 \times 10^{2}$ | 14.33 | - |
| NGC4236 | $1.44 \times 10^{3}$ | 21.26 | $1.92 \times 10^{3}$ | 18.61 | - |
| NGC4254 | $2.51 \times 10^{2}$ | 11.04 | $5.78 \times 10^{2}$ | 9.77 | - |
| NGC4321 | $1.48 \times 10^{2}$ | 12.66 | $2.92 \times 10^{2}$ | 11.91 | - |
| NGC4536 | $1.53 \times 10^{2}$ | 14.29 | $7.73 \times 10^{2}$ | 11.92 | - |
| NGC4559 | $6.18 \times 10^{2}$ | 19.46 | $7.66 \times 10^{2}$ | 17.22 | - |
| NGC4569 | $1.14 \times 10^{2}$ | 6.98 | $1.80 \times 10^{2}$ | 6.83 | - |
| NGC4579 | $1.03 \times 10^{2}$ | 6.98 | $1.82 \times 10^{2}$ | 6.32 | - |
| NGC4594 | $3.22 \times 10^{1}$ | 19.20 | $4.67 \times 10^{1}$ | 18.12 | - |
| NGC4625 | $7.13 \times 10^{2}$ | 21.67 | $7.68 \times 10^{2}$ | 20.86 | - |
| NGC4631 | $3.91 \times 10^{2}$ | 17.10 | $9.49 \times 10^{2}$ | 11.09 | - |
| NGC4725 | $1.79 \times 10^{2}$ | 14.08 | $4.33 \times 10^{2}$ | 10.70 | - |
| NGC4736 | $2.07 \times 10^{2}$ | 13.65 | $7.21 \times 10^{2}$ | 11.50 | - |
| DDO154 | - | - | - | - | 3 |
| NGC4826 | $2.44 \times 10^{2}$ | 9.99 | $3.98 \times 10^{2}$ | 8.85 | - |
| DDO165 | - | - | - | - | 3 |
| NGC5055 | $2.14 \times 10^{2}$ | 13.57 | $4.10 \times 10^{2}$ | 9.95 | - |
| NGC5398 | $\geq 1.12 \times 10^{2}$ | - | $\geq 1.12 \times 10^{2}$ | - | 2 |
| NGC5408 | $5.27 \times 10^{3}$ | 30.20 | $8.74 \times 10^{3}$ | 27.41 | - |
| NGC5457 | $\geq 2.42 \times 10^{2}$ | - | $\geq 2.42 \times 10^{2}$ | - | 2 |
| NGC5474 | $\leq 3.82 \times 10^{2}$ | - | $\leq 4.85 \times 10^{2}$ | - | 2 |
| NGC5713 | $4.42 \times 10^{2}$ | 12.58 | $1.61 \times 10^{3}$ | 8.15 | - |
| NGC5866 | $\leq 2.12 \times 10^{2}$ | - | $\leq 4.68 \times 10^{2}$ | - | 1 |
| NGC6946 | $1.54 \times 10^{2}$ | 9.47 | $3.94 \times 10^{2}$ | 6.83 | - |
| NGC7331 | $2.06 \times 10^{2}$ | 12.54 | $5.61 \times 10^{2}$ | 9.20 | - |
| NGC7793 | $\geq 1.89 \times 10^{2}$ | - | $\geq 1.89 \times 10^{2}$ | - | 2 |
| G11 |  |  |  |  |  |
| M83 | $7.62 \times 10^{2}$ | 9.19 | $8.04 \times 10^{2}$ | 9.09 | - |
| NGC1808 | $2.01 \times 10^{2}$ | 34.57 | $1.16 \times 10^{2}$ | 35.21 | - |
| NGC7552 | $\geq 9.14 \times 10^{1}$ | - | $\geq 9.14 \times 10^{1}$ | - | 2 |
| M82 | $2.70 \times 10^{2}$ | 33.72 | $3.86 \times 10^{2}$ | 33.36 | - |
| NGC1068 | $4.89 \times 10^{1}$ | 36.31 | $2.05 \times 10^{1}$ | 37.50 | - |

Table 9.3. G/D for the DGS, KINGFISH and G11 samples (continued).

| Name | $\mathrm{G} / \mathrm{D}\left(X_{C O, M W}\right)$ | Uncertainty (\%) | $\mathrm{G} / \mathrm{D}\left(X_{C O, Z}\right)$ | Uncertainty (\%) $^{\text {Comments }^{a}}$ |  |
| :--- | :---: | :---: | :---: | :---: | :---: |
| NGC0891 | $3.23 \times 10^{2}$ | 14.96 | $2.37 \times 10^{2}$ | 17.21 | - |
| MGC+02-04-025 | $\geq 2.38 \times 10^{1}$ | - | $\geq 2.38 \times 10^{1}$ | - | 2 |
| NGC7469 | $7.05 \times 10^{1}$ | 19.41 | $6.79 \times 10^{1}$ | 19.43 | - |
| NGC5256 | $\geq 6.86 \times 10^{1}$ | - | $\geq 1.43 \times 10^{2}$ | - | - |
| NGC5953 | $1.36 \times 10^{2}$ | 19.37 | $1.19 \times 10^{2}$ | 19.51 | - |
| M51 | $3.33 \times 10^{1}$ | 22.04 | $4.66 \times 10^{1}$ | 21.93 | - |
| NGC3995 | $\geq 1.50 \times 10^{2}$ | - | $\geq 1.50 \times 10^{2}$ | - | - |
| NGC3994 | $\geq 7.44 \times 10^{1}$ | - | $\geq 7.44 \times 10^{1}$ | - | 2 |
| NGC6052 | $1.63 \times 10^{2}$ | 33.79 | $1.77 \times 10^{2}$ | 33.61 | 2 |
| NGC1222 | $2.01 \times 10^{2}$ | 28.53 | $2.56 \times 10^{2}$ | 27.84 | - |
| NGC7674 | $2.39 \times 10^{2}$ | 31.36 | $1.27 \times 10^{3}$ | 29.75 | - |
| NGC4670 | $2.13 \times 10^{2}$ | 26.05 | $2.68 \times 10^{2}$ | 24.35 | - |

${ }^{a}$ : This column gives the reliability of the point, and works the same way as for the colors in Fig. 9.4.
1 - HI masses are non-detections (blue points in Fig. 9.4).
$2-\mathrm{H}_{2}$ is not detected or not observed for this galaxy (blue points in Fig. 9.4).
3 - The galaxy is not detected beyond $160 \mu \mathrm{~m}$ (red points in Fig. 9.4).
$1+3$ or $2+3$ or $1+2+3$ reflects a combinaison of the previous notes (purple points in Fig. 9.4).
A dash means that we are confident about the three quantities (black points in Fig. 9.4).

## Conclusions

## Summary of activities and results

The general goal of this work was to study the impact of metallicity on the observed ISM properties in galaxies and address the following questions:

- How do the dust properties in galaxies (such as mass, temperature, luminosity, etc.) vary with metallicity? And how do dust model choices influence the overall resulting dust properties?
- In which environments does the submm excess preferentially appear?
- How do the proportions of metals in the gas and dust phases relate to each other? How do the gas-to-dust mass ratios (G/D) evolve as a function of metallicity?

To conduct our study, a multi-wavelength approach has been followed over the whole IR-tosubmm range. The data were interpreted first with a simple modified blackbody model, and second with a more complete semi-empirical dust model allowing a more realistic description of the dust properties.

## Building up the database

I was lucky to start this work about a year after the launch of Herschel when the DGS data was already available for about half of the sample and when members of the PACS and SPIRE ICC had already overcome the major problems in the (complicated) Herschel data reduction. I was thus able to compare several methods for the data reduction and start the photometry on a few DGS galaxies for which the data were already available, before generalising for the whole sample once the observations were complete (in June 2011, nine months after the beginning of my PhD). The major drawback of analysing data from an instrument that is still operating, is the numerous and seemingly never-ending updates of the calibration. The PACS and SPIRE photometry are at the core of my analysis, (especially in the second part of this manuscript), and I did not count the numerous times where I had to redo the maps and photometry for the DGS galaxies to match the most recent update of the map-making technique and of the instrument calibration. It was important to keep track of the calibration: one of the most recent updates had the effect of decreasing the SPIRE flux densities by $\sim 10 \%$, which could have had a non-negligible impact on the results we derive for the dust mass and the submm excess. The other perfect timing that I was very lucky to benefit from was the publication of the KINGFISH Herschel photometry by Dale et al. (2012) at the exact same moment where I was starting to analyse the DGS photometry with modified blackbodies. Including KINGFISH galaxies in my work was thus quite straightforward, and was essential to my work by greatly increasing the metallicity range I was using.

Performing a multi-wavelength approach as done in this work implies a considerable amount of literature searching and data homogenising. This was particularly true for the DGS galaxies, a bit less for the KINGFISH galaxies, as the KINGFISH sample is derived from the Spitzer-SINGS sample,
and most of the IR data for the KINGFISH sample had already been published and homogenised. For the DGS sample however, I searched the literature and databases to find and homogenise the data when available. It also involved treating Spitzer data for the DGS that had been taken in anticipation of the Herschel observations as a set of ancillary data for the DGS. MIPS data had already entirely been reprocessed and published by Bendo et al. (2012). I finalised the treatment of Spitzer data by performing photometry for the IRAC data and treating the IRS spectra to complete our wavelength coverage. In addition to the IR-to-submm data already existing, I wrote several successful proposals to complete our dataset in the submm. I was able to obtain new LABOCA data for three DGS galaxies and conduct the observations on APEX in Chile.

Metallicity, being our main test parameter in this study, required that I ensure that the metallicities were properly and homogeneously determined. Again this had already been done by Moustakas et al. (2010) for the SINGS/KINGFISH galaxies. For the DGS galaxies, I searched the literature to find the required optical line ratios and recomputed the metallicities using the same method and the same strong-line method calibration (Pilyugin \& Thuan 2005) as for the KINGFISH galaxies.

And finally, I also gathered Hi and CO data from the literature for the samples to derive the total gas masses and then G/D for the considered galaxies.

## Study of the global dust properties

We showed that a good wavelength coverage of the peak and of the Rayleigh Jeans slope of the dust SED, as now possible with Herschel, is crucial if one wants to accurately determine the dust properties such as the dust mass and TIR luminosity. A realistic description of dust is also necessary to properly estimate the dust mass. Indeed we found that dust masses estimated using a modified blackbody model are systematically underestimated by a factor of $\sim 1.8$ compared to dust masses estimated with a full dust SED model. This can have important consequences for studies using modified blackbodies to derive dust masses, and for example, go further to estimate a G/D using this underestimated dust mass.

Combining FIR/submm colours with results from the full dust SED model and from the modified blackbody model, we showed that the dust SEDs in dwarf galaxies can have a very peculiar shape: they often have a broader peak of their dust SED, that is shifted to shorter wavelengths compared to the dust SED of more metal-rich environments. This reflects a higher average dust equilibrium temperature ( $\mathrm{T}_{d} \sim 30 \mathrm{~K}$ ) and a broader dust equilibrium temperature distribution ( $\Delta T_{d} \sim 110 \mathrm{~K}$ ) in the low-metallicity galaxies compared to more metal-rich environments ( $\mathrm{T}_{d} \sim 20 \mathrm{~K}$ and $\Delta T_{d} \sim$ 40 K ). This overall warmer dust is due to the harder ISRF interacting with the porous ISM of dwarf galaxies (e.g., Madden et al. 2006). This effect had already been observed for a handful of dwarf galaxies (e.g., Galliano et al. 2003, 2005; Walter et al. 2007; Engelbracht et al. 2008; Galametz et al. 2009) and is now confirmed on a larger sample of dwarf galaxies and quantified. As a consequence, we also found that low-metallicity dwarf galaxies emit more luminosity per unit dust mass over the whole IR range (about seven times more) than more-metal rich environments. This is directly linked to the harder ISRF observed in low-metallicity galaxies along with the low dust abundances.

We also explored dust-to-stellar mass ratios and specific star-formation rates and found that metallicity only has an impact of the second of these two parameters: the dust-to-stellar mass ratios is more or less constant over the metallicity range we are probing, but we found that dwarf galaxies are more efficient in forming stars than their metal-rich counterparts.

## Study of the submm excess

We investigated the presence of an excess at submm wavelengths in our sample with a full dust model for a subsample of 78 galaxies detected at $500 \mu \mathrm{~m}$. We found an excess at $500 \mu \mathrm{~m}$ in $35 \%$ of
the cases and not only in metal-poor galaxies. When including longer submm wavelengths (at 850, 870 or $1200 \mu \mathrm{~m}$ ) we found that about half of the sample presents an excess in the submm. This shows that the submm excess is not uncommon and not only restricted to low-metallicity galaxies. We investigated the presence of the submm excess when using a modified blackbody model with a fixed emissivity, $\beta=2.0$, and found coherent results compared to the submm excess from a full dust SED model. This enabled us to define a method to detect galaxies that have a strong chance of presenting a submm excess from a full dust model, by using Herschel data only and modified blackbodies. We also looked at the influence of metallicity on the excess. We showed that the submm excess appears preferentially in low-metallicity galaxies, but that the intensity of the excess does not depend on metallicity. At low metallicities (for $12+\log (\mathrm{O} / \mathrm{H})<8.0$ ), the mean starlight intensity, which is related to the dust average temperature, discriminates between the galaxies with or without excess: galaxies with low dust temperatures, given their low metallicity ( $\sim 20 \mathrm{~K}$ ), do not show an excess. FIR/submm colours also provide a good empirical tool to detect and predict the appearance of the submm excess. Galaxies showing an excess at $500 \mu \mathrm{~m}$ have high $\frac{F_{160} / F_{250}}{F_{250} / F_{500}}$ ratios ( $\gtrsim 0.37$ ). Galaxies with $\frac{F_{160} / F_{250}}{F_{250} / F_{500}} \gtrsim 0.37$ that do not show an excess at $500 \mu \mathrm{~m}$, either exhibit it at longer wavelength or are low-metallicity galaxies with low dust temperatures given their metallicity. At longer wavelengths, we find that the slope of the submm SED determines whether or not an excess is present: galaxies with $F_{500} / F_{850(870)} \lesssim 5.5$ present an excess for wavelengths $\geq 850 \mu \mathrm{~m}$. We have thus highlighted in this study various observational and modelling criteria to identify galaxies that present an excess. This provides useful tools to determine if a galaxy has an excess when there is limiting data to model it with a full dust SED model.

Additionally we model the IR-to-submm SEDs of the excess galaxies with another dust model with an alternative dust composition: graphite grains are replaced by amorphous carbon grains. We were able to explain the excess in 12 galaxies out of the 37 presenting an excess at submm wavelength with the graphite grains model. These galaxies are mostly moderately metal-poor galaxies and had the weakest excesses with the graphite grains model. We thus believe that the submm excess in these galaxies may be due to the fact that a Galactic grain composition is not appropriate to model the dust emission in these galaxies. This confirms the results of Galliano et al. (2011); Galametz et al. (2013) who showed that a dust composition using amorphous carbon was also more appropriate for the moderately low-metallicity LMC. Nonetheless, a submm excess is still present in 25 galaxies of our sample when modelled with amorphous carbon grains, a sign that other mechanisms are needed to explain the excess in these galaxies.

## Study of the variation of $G / D$ and metallicity

We also studied the relation between G/D and metallicity using more accurate dust masses compared to previous studies of the G/D using Spitzer-only dust masses and/or modified blackbody dust masses. We also probed this parameter over an unprecedented metallicity range ( 2 dex) with a significant fraction of the sample below $12+\log (\mathrm{O} / \mathrm{H})=8.0(\sim 30 \%)$, enabling us to better constrain the behaviour of the G/D at low metallicities. We model the observed trend of the G/D with metallicity using a simple power law (slope of -1 and free) and a broken power law. We found that the G/D does not follow the "reference" trend with metallicity (single power law with a slope of -1 ) over the whole metallicity range. The observed trend is steeper for metallicities lower than $\sim$ 8.0. A large scatter is observed in the G/D values for a given metallicity: in metallicity bins of $\sim 0.1$ dex, the dispersion around the mean value is 0.37 dex on average. On average, the broken power law reproduces best the observed G/D compared to the two power laws (slope of -1 or free) and provides estimates of the G/D that are accurate to a factor of 1.6. Even though more data for galaxies with $12+\log (\mathrm{O} / \mathrm{H}) \leq 7.5$ are necessary to better constrain the observed G/D at extremely low metallicities, this broken power law relation is the best empirical estimate of the observed G/D
for local galaxies we have at our disposal so far. This new empirical prescription can be used for local galaxies to estimate the G/D based on a metallicity value (where this metallicity value must be estimated with the same method than the one we used). Predictions from three chemical evolution models are also compared to the observed G/D values: a simple model from Galliano et al. (2008), a model from Asano et al. (2013) which includes dust growth in the ISM as a source of dust production and continuous star formation; and a model from Zhukovska (2014), more specific to dwarf galaxies, testing episodic star-formation histories and its effects on dust growth and the G/D. The good agreement of observed values of the $G / D$ and its scatter with respect to metallicity with the predicted values of the three tested chemical evolution models, allows us to infer that the scatter in the relation is intrinsic to galactic properties, reflecting the different star-formation histories, dust destruction efficiencies, dust grain size distributions and chemical compositions across the sample.

## Publications of the results

The main results of this work are published or in the process of being published. A first firstauthor paper has been published in Astronomy \& Astrophysics (A\&A 557, A95) presenting the photometry for the DGS sample and the first conclusions I derived on the dust properties with modified blackbodies (i.e., Chapters 4 and 5). A second first-author paper presenting the G/D versus metallicity results (Chapter 9) has been recently accepted by A\&A (acceptance on the 9th of December 2013, arXiv 1312.3442). A third first-author paper is in preparation to present the results obtained on the dust properties and the submm excess with the full dust model, over the whole IR range (i.e., Chapters 6, 7 and 8). Additionally, I am also second author in Madden et al. (2013) where I was in charge of collecting all of the ancillary data for the presentation of the DGS sample. I am also a co-author in ten other papers from the Herschel SAG2 consortium. I was also lucky to have the opportunity of presenting my work at conferences through posters or oral presentations. I was also invited twice to give seminars about my work in other institutes.

Another major input of my work is the wide database of maps/photometry/parameters I put together for the DGS galaxies that should be made available for the community in the next few months.

## Perspectives

This work opens a certain number of perspectives that I would like to explore in future works.

## Zooming into the ISM

Galliano et al. (2011) recently modelled the LMC at different spatial resolutions and found that the lack of resolution could lead to an underestimation of the dust mass of about $50 \%$. Modelling unresolved distant galaxies with the current available models can thus lead to large uncertainties on the dust properties, due to the mixing of the different ISM phases into a single beam: despite its small filling factor, the warm dust phase around the star forming regions can often dominate the emission. To interpret the observations correctly, spatial studies must zoom into the ISM structures to establish what processes drive the physics of these different regions. The wealth of available Herschel data on well resolved local galaxies, coupled with the existing Spitzer and ground-based data, makes it a perfect time to study the effect of varying spatial scales on the interpretation of SEDs of a wide variety of spatially resolved regions of galaxies from hot Hir regions around young star clusters to cold dense molecular clouds. It is possible to relate the observed variations in the dust SEDs and gas properties to the actual physical phenomena (for example the heating of cold dust by the older, more evolved, stellar population), occurring within the ISM of the galaxy (e.g.,

Bendo et al. 2012; Parkin et al. 2012; Aniano et al. 2012; Galametz et al. 2012). The study of a variety of dust environments within the same galaxy (for example the LMC, SMC, M31 or M51, 4 galaxies extensively studied with Herschel) will unveil the dust physical conditions of these regions (intensity of the radiation field, temperature and grain size distributions, abundances, etc.). A few dwarf galaxies from the DGS also show resolved details with Herschel (SF regions, dust embedded young super star clusters...) out to submm wavelengths. With these few test cases we can study how the metallicity affects the impact of spatial resolution on the dust properties (such as the dust mass), allowing the construction of a rich interpretative framework for unresolved, more distant galaxies in the early Universe. These initial studies will undoubtedly be compelling cases to followup with ALMA, for example, to obtain the best spatial view of the structure of the ISM and to see how this affects dust mass determination in galaxies.

## Follow-up on the submm excess

Our study has now, systematically identified galaxies harbouring a submm excess. An obvious follow-up on the submm excess project would be to extend the Herschel observations to longer wavelengths with more APEX/LABOCA or IRAM/GISMO (observing at 2 mm ) observing campaigns, or by using the Planck database following the Early Release Compact Source Catalog and the Planck Catalog of Compact Sources; in cases where there is already a suggestion of submm excess at $500 \mu \mathrm{~m}$ but no observations at longer wavelengths. Using the criteria I defined to find the submm excess in galaxies could also be applied to other Herschel samples, for example the Herschel Reference Survey or the Herschel Virgo Cluster Survey, where there has already been evidence for submm excess (with modBB models, see Grossi et al. 2010) to make a more thorough selection of galaxies with potential submm excess in these samples. This would enrich the first systematic study we conducted on the submm excess at the end of Chapter 8 and see if our results are confirmed on an even larger number of galaxies. Such a study is now possible thanks to the wealth of existing Herschel data.

Application of various explanations proposed for the submm excess, such as spinning dust, and emissivity effects due to an inverse temperature- $\beta$ correlation, on galaxies for which the excess remains after altering the dust composition, is a follow-up project that I would like to develop.

Galliano et al. (2011) found that the submm excess correlates with the diffuse ISM in the LMC. Thus zooming into smaller spatial scales, for example with ALMA in the submm, is a next step to better understand the origin of this excess and the local properties of the ISM that favour the manifestation of the excess. This will be an obvious follow-up to propose for some of the well-justified galaxies already showing a global submm excess. My work shows that the excess is not uncommon even in moderately metal-poor environments. Thus, one would expect the submm excess to also appear in the outer parts of large spiral galaxies where the metallicity can decrease by up to a factor of 10 between the centre and outer parts. Globally those galaxies do not show an excess but they may on local scales, hence highlighting the importance of zooming into the smaller scales of the ISM.

## Constraining the $G / D$ at extremely low metallicities

As a consequence of this work, the behaviour of the $G / D$ is relatively well constrained at low metallicities. However, more extremely-low-metallicity galaxies $(12+\log (\mathrm{O} / \mathrm{H}) \leq 7.5)$ are needed to strengthen our results. This study was based on a sample of rather gas-rich dwarf galaxies. To follow-up on this project we could look for more extremely metal-poor dwarf and/or gas-poor dwarf in other Herschel samples to see how their observed G/D compare with the derived trends and the considered chemical evolution models. There are a number of extremely low-metallicity galaxies that will soon become available in the Herschel archive. A similar treatement of the Herschel data
as presented here, along with the SED modelling, can be carried out to refine the low-metallicity behaviour of the $\mathrm{G} / \mathrm{D}$ below $12+\log (\mathrm{O} / \mathrm{H})=7.5$. Dust and gas masses at this low-metallicity regime will be very important to chemical evolution models which are at the heart of understanding the process of evolution of galaxies.

## Glossary

| 2MASS | 2 Microns All-Sky Survey | HFI | High Frequency Instrument (on Planck) |
| :---: | :---: | :---: | :---: |
| Ac | Amorphous carbon | HIFI | Heterodyne Instrument for the Far Infrared |
| AdOpt | Advanced Optimal extraction (in CASSIS) | HIM | Hot Ionised Medium |
| AGB | Asymptotic Giant Branch | HIPE | Herschel Interactive Processing Environment |
| AGN | Active Galactic Nuclei | HRS | Herschel Reference Survey |
| ALMA | Atacama Large Millimetre/ submillimetre Array | ICC | Instrument Control Centre |
| APEX | Atacama Path finder EXperiment | ICF | Ionisation Correction Factors |
| BCD | Blue Compact Dwarf | IMF | Initial Mass Function |
| BG | Big Grains | IR | InfraRed |
| BoA | Bolometer Array analysis software (for LABOCA) | IRAC | InfraRed Array Camera |
| CASSIS | Cornell AtlaS of Spitzer Infrared spectrograph Sources | IRAM | Institute for Radio Astronomy at Millimetre wavelengths |
| cm | centimetre | IRAS | InfraRed Astronomical Satellite |
| CMB | Cosmic Microwave Background | IRS | InfraRed Spectrograph |
| CNM | Cold Neutral Medium | ISM | InterStellar Medium |
| COBE | COsmic Background Explorer | ISO | Infrared Space Observatory |
| DCD | Disordered Charge Distribution | ISOCAM | ISO CAMera |
| dE | dwarf Elliptical | ISRF | InterStellar Radiation Field |
| DGS | Dwarf Galaxy Survey | JCMT | James Cerk Maxwell Telescope |
| DIRBE | Diffuse InfraRed Background Experiment | KINGFISH | Key Insights on Nearby Galaxies: a Far-Infrared Survey with Herschel |
| dIrr | dwarf Irregular | LABOCA | Large Apex BOlometer CAmera |
| dS | dwarf Spiral | LFI | Low Frequency Instrument (on Planck) |
| dSph | dwarf Spheroidal | LH | Long-High (IRS module) |
| eef | encircled energy fraction | LINER | Low-Ionisation Nuclear Emission-line Regions |
| ESO | European Southern Observatory | LIRG | Luminous InfraRed Galaxy |
| FIR | Far-InfraRed | LL | Long-Low (IRS module) |
| FIRAS | Far InfraRed Absolute Spectrometer | LMC | Large Magellanic Cloud |
| FIRST | Far InfraRed and Submillimetre space Telescope | LW | Long Wave - SCUBA |
| FOV | Field Of View | MAMBO | MAx-planck-Millimetre-BOlometer |
| FUV | Far Ultra-Violet | MIPS | Multiband Imaging Photometer for Spitzer |
| FWHM | Full Width at Half Maximum | MIR | Mid-InfraRed |
| GALEX | GALaxy evolution EXplorer | mm | milimetre |
| G/D | Gas-to-Dust mass ratio | $\operatorname{modBB}$ | modified BlackBody |
| Gr | Graphite | NIR | Near-InfraRed |


| OSO | Onsala Space Observatory | SPIRE | Spectral and Photometric Imaging REceiver <br> PACS |
| :--- | :--- | :--- | :--- |
|  | Photodetector Array Camera <br> and Spectrometer |  |  |
| PSCH | Polycyclic Aromatic Hydrocarbon |  |  |
| PEGASE | Projet d'Étude des GAlaxies | submm | submillimetre |
|  | par Synthèse Évolutive | SW | Short Wave - SCUBA |
| PDR | Photo-Dissociation Region |  |  |
| PSF | Point Spread Function | TDA | Thermal Discrete Approximation |
| RSRF | Relative Spectral Response Function | THINGS | The Hi Nearby Galaxy Survey |
| SCUBA | Submillimetre Common-User | TIR | Total InfraRed |
|  | Bolometer Array | TLS | Two Level System |
| SED | Spectral Energy Distribution |  |  |
| SEST | Swedish-ESO Submillimetre Telescope | UIB | Unidentified Infrared Bands |
| SFR | Star-Formation Rate | UV | Ultra-Luminous InfraRed Galaxy |
| SH | Short-High (IRS module) | VCD | Ultra-Violet |
| SL | Short-Low (IRS module) | VNGS | Very Nearby Galaxy Survey |
| SMART | Spectroscopic Modeling Analysis | VSG | Very Small Grain |
|  | and Reduction Tool (tool within CASSIS) |  |  |
| SMC | Small Magellanic Cloud | WIM | Warm Ionised Medium |
| S/N | Signal-to-Noise ratio | WISE | Wide-field Infrared Survey Explorer |
| SN | SuperNova | WNM | Warm Neutral Medium |
| SNII | Type II Super-Nova | WR | Wolf-Rayet |

## Notations

| $a$ | dust grain radius |
| :---: | :---: |
| $a_{\text {max }}$ | maximum dust grain radius |
| $a_{\text {min }}$ | minimum dust grain radius |
| $a_{t}$ | transition radius between small and big grains, fixed to 10 nm (Chap. 7) |
| $\mathcal{A}$ | telescope area (Chap. 6) |
| $A(\lambda)$ | extinction |
| $A_{u l}$ | Einstein coefficient for spontaneous emission (Chap. 2) |
| $A_{V}$ | V-band extinction |
| $\alpha$ | power-law index of the starlight intensity distribution |
| $\alpha_{0}$ | power-law index of the power-law fit to the observed G/D (Chap. 9) |
| $\alpha_{H}$ | high-metallicity slope of the broken power-law fit to the observed G/D (Chap. 9) |
| $\alpha_{L}$ | low-metallicity slope of the broken power-law fit to the observed G/D (Chap. 9) |
| $\alpha_{S}$ | spectral index (for colour correction) |
| $\alpha_{C O}$ | X-factor in terms of mass surface density |
| $\alpha_{C O, \odot}$ | Galactic $\alpha_{C O}\left(3.8 M_{\odot}\left(\mathrm{K} \mathrm{km} \mathrm{s}^{-1} \mathrm{pc}^{2}\right)^{-1}\right)$ |
| $\alpha_{i j}$ | mass fraction of the $\mathrm{j}^{\text {th }}$ element in the $\mathrm{i}^{\text {th }}$ dust specie (Chap. 7) |
| $\alpha(\lambda)$ | absorption coefficient |
| $\alpha_{\text {sync }}$ | spectral index of the synchrotron spectrum (Chap. 6) |
| $b$ | background level (PACS photometry, Chap. 4) |
| $B_{\nu}(\lambda, T)$ | Planck function |
| $\beta$ | emissivity index |
| $\beta_{o b s}$ | effective emissivity index from a modified blackbody fit with free emissivity index |
| $\beta_{\text {theo }}$ | theoretical emissivity index (from models) |
| $c$ | speed of light |
| C | calibration factor for an instrument (Chap. 6) |
| $C_{a b s}(a, \lambda)$ | dust absorption cross-section |
| $C_{\text {ext }}(a, \lambda)$ | dust extinction cross-section |
| $C_{L L}$ | constant correction factor for Long-Low (Chap. 6) |
| $C_{\text {sca }}(a, \lambda)$ | dust scattering cross-section |
| $C_{S L}(\lambda)$ | wavelength-dependent correction factor for Short-Low (Chap. 6) |
| $C(T)$ | specific heat per unit volume |


| $D_{1}$ | distance |
| :--- | :--- |
| $D_{25}$ | diameter corresponding to $R_{25}$ |
| $D E C$ | Declination |
| $\Delta R(\lambda)$ | error on the relative residual at the wavelength $\lambda$ |
| $\Delta \nu$ | bandwidth of a filter |


| $l$ | length of a given line-of-sight (Chap. 1) |
| :---: | :---: |
| $l_{\nu}^{\text {dust }}$ | SED emitted by a single mass-element of the ISM (Chap. 7) |
| $l_{\nu}^{i}$ | monochromatic luminosities per unit dust mass of the $\mathrm{i}^{\text {th }}$ dust specie, integrated over the size distribution (Chap. 7) |
| $l_{\nu}^{\star}$ | monochromatic luminosities per unit stellar mass (Chap. 7) |
| $L_{C O}$ | CO luminosity |
| $L_{\text {FIR }}$ | far-infrared luminosity |
| $L_{H \alpha}$ | $\mathrm{H} \alpha$ luminosity |
| $L_{H \beta}$ | $\mathrm{H} \beta$ luminosity |
| $L_{\nu}$ | monochromatic luminosity |
| $L_{\nu}^{f f}$ | free-free monochromatic luminosity |
| $L_{\nu}^{\text {model }}$ | synthetic monochromatic luminosity from a given model |
| $L_{\nu}^{\text {observed }}$ | observed monochromatic luminosity |
| $L_{\nu}^{\text {radio }}$ | radio monochromatic luminosity |
| $L_{\nu}^{\star}$ | stellar monochromatic luminosity |
| $L_{\nu}^{\text {sync }}$ | synchrotron monochromatic luminosity |
| $L_{\nu}^{t o t}$ | total monochromatic luminosity |
| $L_{\nu}^{\text {tot, }{ }^{\text {dust }} \text {, }}$ | total dust monochromatic luminosity |
| $L_{\text {TIR }}$ | total infrared luminosity |
| $L_{\text {TIR }}(S)$ | total infrared luminosity estimated only with Spitzer constraints (Chap. 8) |
| $L_{\text {TIR }}(S+H)$ | total infrared luminosity estimated with Spitzer and Herschel constraints (Chap. 8) |
| $L_{\text {tot }}$ | total luminosity of a galaxy |
| $\lambda_{0}$ | designates a reference wavelength |
| $\Lambda$ | cooling rate (Chap. 2) |
| $m(\lambda)$ | complex refractory index (Chap. 1) |
| $\operatorname{mag}(\lambda)$ | apparent magnitude |
| $m_{e}$ | mass of the electron |
| $m_{g}$ | mass of a single dust grain (Chap. 1) |
| $m_{B}$ | B-band apparent magnitude |
| $m_{H}$ | mass of a hydrogen atom |
| $m_{H_{2}}$ | mass of one $\mathrm{H}_{2}$ molecule |
| $m_{j}$ | atomic mass of the $\mathrm{j}^{\text {th }}$ chemical element (Chap. 7) |
| $\operatorname{Mag}(\lambda)$ | absolute magnitude |
| $M_{B}$ | B-band absolute magnitude |
| $M_{d}(S)$ | dust mass estimated only with Spitzer constraints (Chap. 8) |
| $M_{d}(S+H)$ | dust mass estimated with Spitzer and Herschel constraints (Chap. 8) |
| $M_{\text {dust }}$ | dust mass from the full dust SED model |
| $M_{\text {dust, } B \text { B }}$ | dust mass from a modified blackbody |
| $M_{g}$ | total mass of the grains (Chap. 1) |
| $M_{H e}$ | helium mass |
| $M_{i}$ | total mass of the grains of the $\mathrm{i}^{\text {th }}$ dust specie (Chap. 7) |
| $M_{\text {gas }}$ | total gas mass |
| $M_{\text {gas }, \odot}$ | total gas mass estimated with $X_{C O, M W}$ (Chap 9) |
| $M_{\text {gas }, Z}$ | total gas mass estimated with $X_{C O, Z}$ (Chap 9) |
| $M_{H I}$ | atomic gas mass |
| $M_{H I}$ (IRcorr) | atomic gas mass within the dust IR aperture (Chap 9) |
| $M_{\text {HII }}$ | ionised gas mass |
| $M_{H_{2}}$ | molecular gas mass |
| $M_{H_{2}, M W}$ | molecular gas mass estimated with $X_{C O, M W}$ (Chap 9) |
| $M_{H_{2}, Z}$ | molecular gas mass estimated with $X_{C O, Z}$ (Chap 9) |
| $M_{\star}$ | stellar mass |
| $M_{\text {tot }}$ | total mass of a galaxy |
| $M_{V}$ | V-band absolute magnitude |
| $\mu_{\text {gal }}$ | mean atomic weight |
| $\mu_{V}$ | V-band surface brightness |
| $n(a)$ | dust grain size distribution (Chap. 1) |
| $n_{c}$ | critical density of a two-level system (Chap. 2) |
| $n_{d}$ | dust number density (Chap. 1) |
| $n_{e}$ | electron density |
| $n_{l}$ | population density of the lower energy level of a two-level system (Chap. 2) |
| $n_{u}$ | population density of the upper energy level of a two-level system (Chap. 2) |
| $N_{a p}$ | number of pixels within the source aperture (Chap. 4, 6) |
| $N_{b g}$ | number of pixels within the background aperture (Chap. 4, 6) |
| $N_{d}$ | dust column density (Chap. 1) |
| $N_{e}^{S}$ | number of electrons collected by the instrument from the source (Chap. 6) |
| $N_{g}$ | number of dust grains along the line-of-sight (Chap. 1) |
| $N_{H}$ | line-of-sight column density of hydrogen (Chap. 7) |
| $N\left(H_{2}\right)$ | molecular gas column density |
| $N_{j}$ | column density of the $\mathrm{j}^{\text {th }}$ chemical element of a dust specie (Chap. 7) |
| $\nu_{0}$ | designates a reference frequency |

$(O / H)$
$(O / H)$
$\omega$
$\omega_{0}$
$\omega_{p}$
$\Omega$
$\Omega_{m b}$
$\Omega(u, l)$
$P(a, T) \quad$ grain temperature probability distribution
$Q_{a b s}(a, \lambda) \quad$ absorption efficiency
$Q_{e x t}(a, \lambda)$
$Q_{s c a}(a, \lambda)$

## $r_{0}$

$r_{1}$
$r_{2}$
$r_{d u s t}$
$r_{\text {HI }}$
$R_{25}$
$R(500)$
$R(850)$
$R(870)$
$R(1200)$
$R A$
$R(\nu)$
$R_{V}$
$\rho$
$\rho_{i}$
$\rho_{I S M}$
$S F R$
$S_{\lambda}^{i}$
$S$
${ }_{S}^{S} / N(\lambda)$
SSFR
$\sigma_{F_{0}}$
$\sigma_{\text {int }, i}$
$\sigma_{\lambda}$
$\sigma_{m a g}$
$\sigma_{s k y}$
$t$
$T$
$T_{C M B}$
$T_{d}$
$T_{d, \odot}$
$T_{e}$
$T_{e}(O I I I)$
$T_{m b}$
$T_{\text {sub }}$
$\tau(\lambda)$
$\tau_{H}(\lambda)$
$\tau_{S F}$
$\theta$
$u_{\lambda}$
$u n c_{a p}$
uncan
$u n c_{b}$
$u_{n} c_{\text {beam }}$
$u n c_{b g}$
$u n c_{\text {calib }}$
$u_{n} c_{\text {flux }}$
$u n c_{f_{\text {tot }}}$
$u n c_{i n t}$
uncsource
$u n c_{\text {sum }}$
$U$
$<U>$
$U_{\min }$
$U_{\min }(S)$
$U_{\min }(S+H)$
$U_{\max }$
oxygen abundance (or $12+\log (\mathrm{O} / \mathrm{H})$ )
pulsation (Chap. 1)
resonant frequency (Chap. 1)
plasma pulsation (Chap. 1)
solid angle
main beam solid angle for radio astronomy (Chap. 6)
collision strength for upward collisions (Chap. 2)
scattering efficiency
source aperture radius (PACS photometry, Chap. 4)
inner background annulus radius (PACS photometry, Chap. 4)
outer background annulus radius (PACS photometry, Chap. 4)
IR aperture dust radius (Chap 9)
HI radius (Chap 9)
residual at $500 \mu \mathrm{~m}$ relative to the model
residual at $850 \mu \mathrm{~m}$ relative to the model
residual at $870 \mu \mathrm{~m}$ relative to the model
residual at $1200 \mu \mathrm{~m}$ relative to the model
Right Ascension
response function of a given band
extinction curve coefficient, ratio of the total-to-selective extinction
dust material mass density (Chap. 1)
mass density of the $\mathrm{i}^{\text {th }}$ dust component (Chap. 7)
cloud mass density (Chap. 1)
star-formation rate
source function for the $\mathrm{i}^{\text {th }}$ dust component (Chap. 7) specific star-formation rate
uncertainty on the zero-magnitude flux value (Chap. 6)
final uncertainty on the flux density at the wavelength $\lambda$
error on the magnitude (Chap. 6)
integration time (Chap. 6)
temperature of the dust from a modified blackbody
temperature of the CMB radiation (2.73 K, Mather et al. 1994)
dust equilibrium temperature
electron temperature
main beam temperature (Chap. 6)
sublimation temperature for the grain (Chap. 7)
optical depth
extinction per H column density (Chap. 7)
star formation timescale (Chap. 9)
FWHM of the main beam for radio astronomy (CO) (Chap. 6)
energy density
uncertainty on the background level (PACS photometry, Chap. 4)
intensity of the stellar radiation field
mass-averaged starlight intensity
minimum intensity of the stellar radiation field maximum intensity of the stellar radiation field
solar neighbourhood oxygen abundance ( $4.90 \times 10^{-4}$ (or 8.69), Asplund et al. 2009)
radius at which the surface brightness falls to a level of $25 \mathrm{mag} / \operatorname{arcsec}^{2}$
signal-to-noise ratio of the excess at a given wavelength (or intensity of the excess)
intrinsic uncertainty coming from the data reduction for one pixel (Chap. 4)
standard deviation of all pixels in the background aperture (Chap. 4, 6)
dust equilibrium temperature of the diffuse Galactic dust (17.5 K, Boulanger et al. 1996)
electron temperature determined from the [OiII] $\lambda 4363$ line (Chap. 3)
uncertainty on the measured flux in the source aperture (PACS photometry, Chap. 4) uncertainty on the measured flux in the background annulus (PACS photometry, Chap. 4) uncertainty from the SPIRE beam area (SPIRE photometry, Chap. 4)
uncertainty from the background determination (SPIRE photometry, Chap. 4) systematic calibration uncertainty on the measured total flux density (Chap. 4) non-systematic uncertainty on the measured total flux density (SPIRE photometry, Chap. 4) non-systematic uncertainty on the measured total flux density (PACS photometry, Chap. 4) intrinsic uncertainty coming from the data reduction for the total flux density (Chap. 4) uncertainty on the measured flux in the source aperture (SPIRE photometry, Chap. 4) uncertainty coming from summing pixels (PACS photometry, Chap. 4)
minimum intensity of the stellar radiation field estimated only with Spitzer constraints
minimum intensity of the stellar radiation field estimated with Spitzer and Herschel constraints

| $W_{i}$ | dilution factors for the optical components of $J_{\lambda}^{\text {Gal }}$ (Chap. 7) |
| :---: | :---: |
| $x=2 \pi a / \lambda$ | size parameter (Chap. 1) |
| $X$ | mass fraction of hydrogen |
| $X_{\odot}$ | mass fraction of hydrogen in the solar neighbourhood (0.7154, Asplund et al. 2009) |
| $X_{C O}$ | conversion factor between CO and $\mathrm{H}_{2}$, so called "X-factor" |
| $X_{C O, M W}$ | Galactic conversion factor between CO and $\mathrm{H}_{2}\left(2.0 \times 10^{20} \mathrm{~cm}^{-2}\left(\mathrm{~K} \mathrm{~km} \mathrm{~s}^{-1}\right)^{-1}\right.$, Ackermann et al. 2011) |
| $X_{C O, Z}$ | metallicity-dependent $X_{C O}$ factor, as in Schruba et al. (2012) |
| $X_{P}$ | primordial mass fraction of hydrogen (0.76, Pagel 1997) |
| Y | mass fraction of helium |
| $Y_{\odot}$ | mass fraction of helium in the solar neighbourhood (0.2703, Asplund et al. 2009) |
| $Y_{P}$ | primordial mass fraction of helium (0.24, Pagel 1997) |
| $Z$ | metallicity (mass fraction of metals) |
| $Z_{\odot}$ | solar neighbourhood metallicity (0.0142, Asplund et al. 2009) |
| $Z_{P}$ | primordial metallicity (0 by definition) |

## Appendix A

Observing Logs

Table A.1. DGS Herschel Observing Log.

| Source | Obs ID | Description | Mapping mode | Duration (sec) | Date of observation |
| :---: | :---: | :---: | :---: | :---: | :---: |
| Haro 11 | 1342210636 | PACS 70-1 | Scan map | 141 | December 2010 |
|  | 1342210637 | PACS 70-2 | Scan map | 153 | December 2010 |
|  | 1342197713 | PACS 100-1 | Scan map | 65 | June 2010 |
|  | 1342197714 | PACS 100-2 | Scan map | 65 | June 2010 |
|  | 1342199386 | SPIRE | Small map | 307 | June 2010 |
| Haro 2 | 1342196745 | PACS 70-1 | Scan map | 262 | May 2010 |
|  | 1342196746 | PACS 70-2 | Scan map | 262 | May 2010 |
|  | 1342198283 | PACS 100-1 | Scan map | 262 | June 2010 |
|  | 1342198284 | PACS 100-2 | Scan map | 262 | June 2010 |
|  | 1342195686 | SPIRE | Large map | 639 | April 2010 |
| Haro 3 | 1342195618 | PACS 70-1 | Scan map | 71 | April 2010 |
|  | 1342195619 | PACS 70-2 | Scan map | 71 | April 2010 |
|  | 1342195616 | PACS 100-1 | Scan map | 71 | April 2010 |
|  | 1342195617 | PACS 100-2 | Scan map | 71 | April 2010 |
|  | 1342207035 | SPIRE | Small map | 307 | October 2010 |
| He 2-10 | 1342196121 | PACS 70-1 | Scan map | 71 | May 2010 |
|  | 1342196122 | PACS 70-2 | Scan map | 71 | May 2010 |
|  | 1342196119 | PACS 100-1 | Scan map | 71 | May 2010 |
|  | 1342196120 | PACS 100-2 | Scan map | 71 | May 2010 |
|  | 1342196888 | SPIRE | Large map | 241 | May 2010 |
| HS 0017+1055 | 1342199634 | PACS 70-1 | Scan map | 220 | July 2010 |
|  | 1342199635 | PACS 70-2 | Scan map | 220 | July 2010 |
|  | 1342199636 | PACS 100-1 | Scan map | 220 | July 2010 |
|  | 1342199637 | PACS 100-2 | Scan map | 220 | July 2010 |
|  | 1342201378 | SPIRE | Small map | 583 | July 2010 |
| HS 0052+2536 | 1342213514 | PACS 70-1 | Scan map | 220 | January 2011 |
|  | 1342213515 | PACS 70-2 | Scan map | 220 | January 2011 |
|  | 1342213516 | PACS 100-1 | Scan map | 220 | January 2011 |
|  | 1342213517 | PACS 100-2 | Scan map | 220 | January 2011 |
|  | 1342201326 | SPIRE | Small map | 583 | July 2010 |
| HS 0822+3542 | 1342209081 | PACS 70-1 | Scan map | 212 | November 2010 |
|  | 1342209082 | PACS 70-2 | Scan map | 212 | November 2010 |
|  | 1342209083 | PACS 100-1 | Scan map | 212 | November 2010 |
|  | 1342209084 | PACS 100-2 | Scan map | 212 | November 2010 |
| HS 1222+3741 | 1342211380 | PACS 70-1 | Scan map | 276 | December 2010 |
|  | 1342211381 | PACS 70-2 | Scan map | 276 | December 2010 |
|  | 1342211382 | PACS 100-1 | Scan map | 276 | December 2010 |
|  | 1342211383 | PACS 100-2 | Scan map | 276 | December 2010 |
| HS 1236+3937 | 1342211434 | PACS 70-1 | Scan map | 220 | December 2010 |
|  | 1342211435 | PACS 70-2 | Scan map | 220 | December 2010 |
|  | 1342197779 | PACS 100-1 | Scan map | 220 | June 2010 |
|  | 1342197780 | PACS 100-2 | Scan map | 220 | June 2010 |
|  | 1342199341 | SPIRE | Small map | 307 | June 2010 |
| HS 1304+3529 | 1342211289 | PACS 70-1 | Scan map | 220 | December 2010 |
|  | 1342211290 | PACS 70-2 | Scan map | 220 | December 2010 |
|  | 1342198475 | PACS 100-1 | Scan map | 164 | June 2010 |
|  | 1342198476 | PACS 100-2 | Scan map | 164 | June 2010 |
|  | 1342201224 | SPIRE | Small map | 307 | July 2010 |
| HS 1319+3224 | 1342211392 | PACS 70-1 | Scan map | 276 | December 2010 |
|  | 1342211393 | PACS 70-2 | Scan map | 276 | December 2010 |
|  | 1342211394 | PACS 100-1 | Scan map | 276 | December 2010 |
|  | 1342211395 | PACS 100-2 | Scan map | 276 | December 2010 |
| HS 1330+3651 | 1342211388 | PACS 70-1 | Scan map | 276 | December 2010 |
|  | 1342211389 | PACS 70-2 | Scan map | 276 | December 2010 |
|  | 1342211390 | PACS 100-1 | Scan map | 276 | December 2010 |
|  | 1342211391 | PACS 100-2 | Scan map | 276 | December 2010 |

Table A.1. DGS Herschel Observing Log.

| Source | Obs ID | Description | Mapping mode | Duration (sec) | Date of observation |
| :---: | :---: | :---: | :---: | :---: | :---: |
| HS 1442+4250 | 1342210958 | PACS 70-1 | Scan map | 317 | December 2010 |
|  | 1342210959 | PACS 70-2 | Scan map | 317 | December 2010 |
|  | 1342210960 | PACS 100-1 | Scan map | 317 | December 2010 |
|  | 1342210961 | PACS 100-2 | Scan map | 317 | December 2010 |
|  | 1342203597 | SPIRE | Small map | 307 | August 2010 |
| HS 2352+2733 | 1342212062 | PACS 70-1 | Scan map | 276 | December 2010 |
|  | 1342212063 | PACS 70-2 | Scan map | 276 | December 2010 |
|  | 1342212064 | PACS 100-1 | Scan map | 276 | December 2010 |
|  | 1342212065 | PACS 100-2 | Scan map | 276 | December 2010 |
|  | 1342201375 | SPIRE | Small map | 307 | July 2010 |
| I Zw 18 | 1342209354 | PACS 70-1 | Scan map | 276 | November 2010 |
|  | 1342209355 | PACS 70-2 | Scan map | 276 | November 2010 |
|  | 1342209356 | PACS 100-1 | Scan map | 276 | November 2010 |
|  | 1342209357 | PACS 100-2 | Scan map | 276 | November 2010 |
|  | 1342187175 | SPIRE | Small map | 2676 | November 2009 |
| IC 10 | 1342216141 | PACS 70-1 | Scan map | 3061 | March 2011 |
|  | 1342216142 | PACS 70-2 | Scan map | 3061 | March 2011 |
|  | 1342216143 | PACS 100-1 | Scan map | 3061 | March 2011 |
|  | 1342216144 | PACS 100-2 | Scan map | 3061 | March 2011 |
|  | 1342201446 | SPIRE | Large map | 2286 | July 2010 |
| II Zw 40 | 1342204313 | PACS 70-1 | Scan map | 153 | September 2010 |
|  | 1342204314 | PACS 70-2 | Scan map | 153 | September 2010 |
|  | 1342204315 | PACS 100-1 | Scan map | 153 | September 2010 |
|  | 1342204316 | PACS 100-2 | Scan map | 153 | September 2010 |
|  | 1342186116 | SPIRE | Large map | 1055 | October 2009 |
| Mrk 1089 | 1342193121 | PACS 70-1 | Scan map | 65 | March 2010 |
|  | 1342193122 | PACS 70-2 | Scan map | 65 | March 2010 |
|  | 1342193172 | PACS 100-1 | Scan map | 65 | April 2010 |
|  | 1342193173 | PACS 100-2 | Scan map | 65 | April 2010 |
|  | 1342192102 | SPIRE | Small map | 193 | March 2010 |
| Mrk 1450 | 1342195596 | PACS 70-1 | Scan map | 182 | April 2010 |
|  | 1342195597 | PACS 70-2 | Scan map | 182 | April 2010 |
|  | 1342195620 | PACS 100-1 | Scan map | 182 | April 2010 |
|  | 1342195621 | PACS 100-2 | Scan map | 182 | April 2010 |
|  | 1342199347 | SPIRE | Small map | 583 | June 2010 |
| Mrk 153 | 1342195827 | PACS 70-1 | Scan map | 108 | May 2010 |
|  | 1342195828 | PACS 70-2 | Scan map | 108 | May 2010 |
|  | 1342197018 | PACS 100-1 | Scan map | 108 | May 2010 |
|  | 1342197019 | PACS 100-2 | Scan map | 108 | May 2010 |
|  | 1342206187 | SPIRE | Small map | 583 | October 2010 |
| Mrk 209 | 1342195850 | PACS 70-1 | Scan map | 65 | May 2010 |
|  | 1342195851 | PACS 70-2 | Scan map | 65 | May 2010 |
|  | 1342197707 | PACS 100-1 | Scan map | 65 | May 2010 |
|  | 1342197708 | PACS 100-2 | Scan map | 65 | May 2010 |
|  | 1342199350 | SPIRE | Small map | 583 | June 2010 |
| Mrk 930 | 1342212066 | PACS 70-1 | Scan map | 317 | December 2010 |
|  | 1342212067 | PACS 70-2 | Scan map | 317 | December 2010 |
|  | 1342212068 | PACS 100-1 | Scan map | 317 | December 2010 |
|  | 1342212069 | PACS 100-2 | Scan map | 317 | December 2010 |
|  | 1342201374 | SPIRE | Small map | 583 | July 2010 |
| NGC 1140 | 1342212756 | PACS 70-1 | Scan map | 909 | January 2011 |
|  | 1342212757 | PACS 70-2 | Scan map | 909 | January 2011 |
|  | 1342212758 | PACS 100-1 | Scan map | 909 | January 2011 |
|  | 1342212759 | PACS 100-2 | Scan map | 909 | January 2011 |
|  | 1342201415 | SPIRE | Large map | 2047 | July 2010 |

Table A.1. DGS Herschel Observing Log.

| Source | Obs ID | Description | Mapping mode | Duration (sec) | Date of observation |
| :---: | :---: | :---: | :---: | :---: | :---: |
| NGC 1569 | 1342206046 | PACS 70-1 | Scan map | 1280 | October 2010 |
|  | 1342206047 | PACS 70-2 | Scan map | 1280 | October 2010 |
|  | 1342206048 | PACS 100-1 | Scan map | 1280 | October 2010 |
|  | 1342206049 | PACS 100-2 | Scan map | 1732 | October 2010 |
|  | 1342193013 | SPIRE | Large map | 529 | March 2010 |
| NGC 1705 | 1342186270 | PACS 70-1 | Scan map | 376 | October 2009 |
|  | 1342186271 | PACS 70-2 | Scan map | 376 | October 2009 |
|  | 1342186300 | PACS 100-1 | Scan map | 376 | October 2009 |
|  | 1342186301 | PACS 100-2 | Scan map | 376 | October 2009 |
|  | 1342186114 | SPIRE | Large map | 2579 | October 2009 |
| NGC 2366 | 1342220565 | PACS 70-1 | Scan map | 1141 | May 2011 |
|  | 1342220566 | PACS 70-2 | Scan map | 1141 | May 2011 |
|  | 1342220567 | PACS 100-1 | Scan map | 1141 | May 2011 |
|  | 1342220568 | PACS 100-2 | Scan map | 1141 | May 2011 |
|  | 1342193014 | SPIRE | Large map | 2047 | March 2010 |
| NGC 4214 | 1342211803 | PACS 70-1 | Scan map | 2792 | December 2010 |
|  | 1342211804 | PACS 70-2 | Scan map | 2792 | December 2010 |
|  | 1342211805 | PACS 100-1 | Scan map | 2792 | December 2010 |
|  | 1342211806 | PACS 100-2 | Scan map | 2792 | December 2010 |
|  | 1342199342 | SPIRE | Large map | 1661 | June 2010 |
| NGC 4449 | 1342221125 | PACS 70-1 | Scan map | 3497 | May 2011 |
|  | 1342221126 | PACS 70-2 | Scan map | 3497 | May 2011 |
|  | 1342221127 | PACS 100-1 | Scan map | 3497 | May 2011 |
|  | 1342221128 | PACS 100-2 | Scan map | 3497 | May 2011 |
|  | 1342198243 | SPIRE | Large map | 1035 | June 2010 |
| NGC 4861 | 1342211384 | PACS 70-1 | Scan map | 685 | December 2010 |
|  | 1342211385 | PACS 70-2 | Scan map | 685 | December 2010 |
|  | 1342211386 | PACS 100-1 | Scan map | 685 | December 2010 |
|  | 1342211387 | PACS 100-2 | Scan map | 685 | December 2010 |
|  | 1342199340 | SPIRE | Large map | 639 | June 2010 |
| NGC 5253 | 1342202376 | PACS 70-1 | Scan map | 1141 | August 2010 |
|  | 1342202377 | PACS 70-2 | Scan map | 1141 | August 2010 |
|  | 1342202378 | PACS 100-1 | Scan map | 1141 | August 2010 |
|  | 1342202379 | PACS 100-2 | Scan map | 1141 | August 2010 |
|  | 1342203078 | SPIRE | Large map | 1035 | August 2010 |
| NGC 625 | 1342213260 | PACS 70-1 | Scan map | 1280 | January 2011 |
|  | 1342213261 | PACS 70-2 | Scan map | 1280 | January 2011 |
|  | 1342213262 | PACS 100-1 | Scan map | 1280 | January 2011 |
|  | 1342213263 | PACS 100-2 | Scan map | 1280 | January 2011 |
|  | 1342195943 | SPIRE | Large map | 1618 | May 2010 |
| NGC 6822 | 1342186167 | PACS 70-1 | Scan map | 2554 | October 2009 |
|  | 1342186168 | PACS 70-2 | Scan map | 2554 | October 2009 |
|  | 1342186169 | PACS 100-1 | Scan map | 2554 | October 2009 |
|  | 1342186170 | PACS 100-2 | Scan map | 2554 | October 2009 |
|  | 1342185533 | SPIRE | Large map | 2579 | October 2009 |
| Pox 186 | 1342213618 | PACS 70-1 | Scan map | 276 | February 2011 |
|  | 1342213619 | PACS 70-2 | Scan map | 276 | February 2011 |
|  | 1342213620 | PACS 100-1 | Scan map | 276 | February 2011 |
|  | 1342213621 | PACS 100-2 | Scan map | 276 | February 2011 |
|  | 1342201238 | SPIRE | Small map | 583 | July 2010 |
| SBS 0335-052 | 1342202302 | PACS 70-1 | Scan map | 276 | August 2010 |
|  | 1342202303 | PACS 70-2 | Scan map | 276 | August 2010 |
|  | 1342202304 | PACS 100-1 | Scan map | 276 | August 2010 |
|  | 1342202305 | PACS 100-2 | Scan map | 276 | August 2010 |
|  | 1342203627 | SPIRE | Small map | 583 | August 2010 |

Table A.1. DGS Herschel Observing Log.

| Source | Obs ID | Description | Mapping mode | Duration (sec) | Date of observation |
| :---: | :---: | :---: | :---: | :---: | :---: |
| SBS 1159+545 | 1342209445 | PACS 70-1 | Scan map | 220 | November 2010 |
|  | 1342209446 | PACS 70-2 | Scan map | 220 | November 2010 |
|  | 1342196852 | PACS 100-1 | Scan map | 276 | May 2010 |
|  | 1342196853 | PACS 100-2 | Scan map | 276 | May 2010 |
|  | 1342199348 | SPIRE | Small map | 583 | June 2010 |
| SBS 1211+540 | 1342197777 | PACS 70-1 | Scan map | 900 | June 2010 |
|  | 1342197778 | PACS 70-2 | Scan map | 900 | June 2010 |
|  | 1342198182 | PACS 100-1 | Scan map | 900 | June 2010 |
|  | 1342199349 | PACS 100-2 | Scan map | 900 | June 2010 |
|  | 1342198181 | SPIRE | Small map | 583 | June 2010 |
| SBS 1249+493 | 1342199123 | PACS 70-1 | Scan map | 220 | June 2010 |
|  | 1342199124 | PACS 70-2 | Scan map | 220 | June 2010 |
|  | 1342199125 | PACS 100-1 | Scan map | 220 | June 2010 |
|  | 1342199126 | PACS 100-2 | Scan map | 220 | June 2010 |
|  | 1342199351 | SPIRE | Small map | 583 | June 2010 |
| SBS 1415+437 | 1342210592 | PACS 70-1 | Scan map | 317 | November 2010 |
|  | 1342210593 | PACS 70-2 | Scan map | 317 | November 2010 |
|  | 1342210594 | PACS 100-1 | Scan map | 317 | November 2010 |
|  | 1342210595 | PACS 100-2 | Scan map | 317 | November 2010 |
| SBS 1533+574 | 1342193505 | PACS 70-1 | Scan map | 153 | April 2010 |
|  | 1342193506 | PACS 70-2 | Scan map | 153 | April 2010 |
|  | 1342193066 | PACS 100-1 | Scan map | 153 | April 2010 |
|  | 1342193067 | PACS 100-2 | Scan map | 153 | April 2010 |
|  | 1342203600 | SPIRE | Small map | 583 | April 2010 |
| Tol 0618-402 | 1342196111 | PACS 70-1 | Scan map | 558 | May 2010 |
|  | 1342196112 | PACS 70-2 | Scan map | 558 | May 2010 |
|  | 1342196113 | PACS 100-1 | Scan map | 558 | May 2010 |
|  | 1342196114 | PACS 100-2 | Scan map | 558 | May 2010 |
| Tol 1214-277 | 1342222748 | PACS 70-1 | Scan map | 671 | June 2011 |
|  | 1342222749 | PACS 70-2 | Scan map | 671 | June 2011 |
|  | 1342222750 | PACS 100-1 | Scan map | 671 | June 2011 |
|  | 1342222751 | PACS 100-2 | Scan map | 671 | June 2011 |
|  | 1342201264 | SPIRE | Small map | 583 | July 2010 |
| UGC 4483 | 1342196016 | PACS 70-1 | Scan map | 141 | May 2010 |
|  | 1342196017 | PACS 70-2 | Scan map | 141 | May 2010 |
|  | 1342196018 | PACS 100-1 | Scan map | 141 | May 2010 |
|  | 1342196019 | PACS 100-2 | Scan map | 141 | May 2010 |
|  | 1342206192 | SPIRE | Small map | 583 | October 2010 |
| UGCA 20 | 1342213565 | PACS 70-1 | Scan map | 293 | January 2011 |
|  | 1342213566 | PACS 70-2 | Scan map | 293 | January 2011 |
|  | 1342213567 | PACS 100-1 | Scan map | 300 | January 2011 |
|  | 1342213568 | PACS 100-2 | Scan map | 293 | January 2011 |
| UM 133 | 1342213553 | PACS 70-1 | Scan map | 293 | January 2011 |
|  | 1342213554 | PACS 70-2 | Scan map | 293 | January 2011 |
|  | 1342213555 | PACS 100-1 | Scan map | 293 | January 2011 |
|  | 1342213556 | PACS 100-2 | Scan map | 293 | January 2011 |
|  | 1342201323 | SPIRE | Small map | 445 | July 2010 |
| UM 311 | 1342213237 | PACS 70-1 | Scan map | 583 | January 2011 |
|  | 1342213238 | PACS 70-2 | Scan map | 583 | January 2011 |
|  | 1342213239 | PACS 100-1 | Scan map | 583 | January 2011 |
|  | 1342213240 | PACS 100-2 | Scan map | 583 | January 2011 |
|  | 1342201321 | SPIRE | Large map | 658 | July 2010 |
| UM 448 | 1342198595 | PACS 70-1 | Scan map | 71 | June 2010 |
|  | 1342198596 | PACS 70-2 | Scan map | 71 | June 2010 |
|  | 1342199769 | SPIRE | Small map | 307 | July 2010 |

Table A.1. DGS Herschel Observing Log.

| Source | Obs ID | Description | Mapping mode | Duration (sec) | Date of observation |
| :--- | :---: | :---: | :---: | :---: | :---: |
| UM 461 | 1342211793 | PACS 70-1 | Scan map | 317 | December 2010 |
|  | 1342211794 | PACS 70-2 | Scan map | 317 | December 2010 |
|  | 1342211795 | PACS 100-1 | Scan map | 317 | December 2010 |
|  | 1342211796 | PACS 100-2 | Scan map | 317 | December 2010 |
|  | 1342200110 | SPIRE | Small map | 307 | July 2010 |
| VII Zw 403 | 1342195400 | PACS 70-1 | Scan map | 71 | April 2010 |
|  | 1342195401 | PACS 70-2 | Scan map | 71 | April 2010 |
|  | 1342195478 | PACS 100-1 | Scan map | 71 | April 2010 |
|  | 1342195479 | PACS 100-2 | Scan map | 71 | April 2010 |
|  | 1342199363 | SPIRE | Small map | 583 | July 2010 |

Table A.2. DGS Spitzer/IRAC and IRS Observing Log.

|  | IRAC | IRS |  |
| :---: | :---: | :---: | :---: |
| Sources | AOR key | AOR key | Extraction |
| Haro 11 | 4326400 | 9007104 | Optimal |
| Haro 2 | 5539840 | 9489920 | Map |
| Haro 3 | 11180288 | 12556288 | Tapered |
| He 2-10 | 4329472 | 4340480 | Tapered |
| HS 0017+1055 | 26387200 | 26393344 | Optimal |
| HS 0052+2536 | 26387456 | 17463040 | Optimal |
| HS $0822+3542$ | 4328960 | 1763808 | Optimal |
| HS 1222+3741 | 17564928 | 26393600 | Optimal |
| HS 1236+3937 | 26387712 | 26393856 | Optimal |
| HS 1304+3529 | 26387968 | 26394112 | Optimal |
| HS 1319+3224 | 26388480 | 26394624 | Optimal |
| HS 1330+3651 | 26388736 | 26394880 | Optimal |
| HS 1442+4250 | 10388480 | 12562944 | Optimal |
| HS $2352+2733^{a}$ | 26388992 | 26395136 | - |
| I Zw 18 | 4330752 | 16205568 | Optimal |
| IC 10 | 4424960 | 26396672 | Map |
| II Zw 40 | 4327936 | 9007616 | Optimal |
| Mrk 1089 | 11250432 | 26395392 | Tapered |
| Mrk 1450 | 4334336 | 16206080 (SL), 9011712 (LL) | Optimal |
| Mrk 153 | 4333056 | 4342272 | Optimal |
| Mrk 209 | 22556672 | 12557568 | Optimal |
| Mrk 930 | 4338944 | 4344320 | Tapered |
| NGC 1140 | 4327168 | 4830976 | Tapered |
| NGC 1569 | 4434944 | 3856640 | Tapered |
| NGC 1705 | 5535744 | 9513216 | Map |
| NGC 2366 | 4436480 | 21920768 | Map |
| NGC $4214{ }^{\text {b }}$ | 4457984 | - | - |
| NGC 4449 | 4467456 | 26396928 | Map |
| NGC 4861 ${ }^{\text {c }}$ | 4337408 | - | - |
| NGC 5253 | 4386048 | 4386304 | Map |
| NGC 625 ${ }^{\text {c }}$ | 22520064 | - | - |
| NGC $6822^{\text {b,c }}$ | 5507072 | - | - |
| Pox 186 | 26389248 | 12629760 | Optimal |
| SBS 0335-052 | 4327424 | 8986880 | Optimal |
| SBS 1159+545 | 26389504 | 9008896 | Optimal |
| SBS 1211+540 | 26389760 | 26395392 | Optimal |
| SBS 1249+493 | 26390016 | 26395904 | Optimal |
| SBS 1415+437 | 10392832 | 12562432 (SL), 8990464 (LL) | Optimal |
| SBS 1533+574 | 17563904 | 8996352 | Optimal |
| Tol 0618-402 | 4328448 | 8090624 | Optimal |
| Tol 1214-277 | 4336384 | 9008128 | Optimal |
| UGC 4483 | 4329728 | 26396160 | Optimal |
| UGCA $20^{a}$ | 26390272 | 26396416 | - |
| UM 133 | 26390528 | 21922304 | Map |
| UM $311^{\text {b }}$ | 10392576 | - | - |
| UM 448 | 4334592 | 4342784 | Tapered |
| UM 461 | 4335104 | 16204032 (SL), 9006336 (LL) | Optimal |
| VII Zw 403 | 4334080 | 9005824 | Tapered |

## Notes :

${ }^{a}$ : For these galaxies, the IRS slits are not centered on the source position and thus we do not present any IRS spectrum.
${ }^{b}$ : For these galaxies, only local pointings have been done and we cannot present an IRS spectrum for the total galaxy.
${ }^{c}$ : Only high resolution IRS spectrum (SH and/or LH) is available.

## Appendix B

The DGS galaxies

RA: 00h36m52.7s $\quad$ DEC : -33d33m17s
D $=92.1 \mathrm{Mpc} \quad 12+\log (0 / \mathrm{H})=8.36$
Morph Type : HIIgalaxy
Other names : ES0350, IG038
Haro11 in the Dwarf Galaxy Survey :


## Haro 2



| RA: 10h32m32.0s | DEC : 54d24m02s |
| :--- | :--- |
| D $=21.7 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=8.23$ |
| Morph Type : Im-pec, WR, HIIgalaxy |  |
| Other names : Mrk33, SBS1029+546, UGC05720 |  |





| RA : 10h45m22.4s | DEC : 57 d 57 m 37 s |
| :--- | :--- |
| $\mathrm{D}=19.3 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=8.28$ |
| Morph Type : BCD, HIIgalaxy |  |
| Other names : Mrk35, NGC3353, SBS1042+562, UGC0586 |  |

Haro3 in the Dwarf Galaxy Survey :


He 2-10

RA: 08h36m15.1s DEC: -26d24m34s
$D=8.7 \mathrm{Mpc} \quad 12+\log (0 / H)=8.43$
Morph Type : IO-pec, starburst
Other names : N/A


## HS 0017+1055



## HS 0052+2536



| RA : 00h20m21.4s | DEC : 11d12m21s |
| :--- | :--- |
| D $=79.1 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=7.63$ |
| Morph Type : BCG |  |
| Other names : N/A |  |



## HS 0822+3542


RA:08h25m55.5s $\quad$ DEC : 35d32m32s
D $=11.0 \mathrm{Mpc} \quad 12+\log (0 / \mathrm{H})=7.32$
Morph Type : BCG, iL,C, BCD
Other names : SDSSJ0825 + 3542


## HS 1222+3741





HS 1304+3529


## HS 1319+3224




HS1330+3651 in the Dwarf Galaxy Survey :


HS 1442+4250


| RA: 14 h 44 m 12.8 s | DEC $: 42 \mathrm{~d} 37 \mathrm{~m} 44 \mathrm{~s}$ |
| :--- | :--- |
| $\mathrm{D}=14.4 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=7.60$ |
| Morph Type : Pec, BCG |  |
| Other names : UGC09497 |  |



| RA : 23h54m56.7s | DEC $: 27 \mathrm{~d} 49 \mathrm{~m} 59 \mathrm{~s}$ |
| :--- | :--- |
| D $=116.7 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=8.40$ |
| Morph Type : N/A |  |
| Other names : N/A |  |






RA: 05h55m42.6s DEC: 3d23m32s
$\mathrm{D}=12.1 \mathrm{Mpc} \quad 12+\log (0 / H)=8.23$
Morph Type : BCD, SBc, merger, HIIgalaxy
Other names: UGCA116

IIZw40 in the Dwarf Galaxy Survey :




| RA: 00h20m17.3s | DEC $: 59 \mathrm{~d} 18 \mathrm{~m} 14 \mathrm{~s}$ |
| :--- | :--- |
| $\mathrm{D}=\quad 0.7 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=8.17$ |
| Morph Type: Ibm, dIrr, BCD |  |
| Other names : UGC00192 |  |

IC10 in the Dwarf Galaxy Survey :


Mrk 1089




Mrk 1450


| RA : 10h49m05.0s | DEC : 52d20m08s |
| :--- | :--- |
| D $=40.3 \mathrm{Mpc}$ | $12+\log (0 / H)=7.86$ |
| Morph Type : Scp |  |
| Other names : SBS1046+526, UGCA219 |  |




RA: 12h26m15.9s DEC: 48d29m37s
$\mathrm{D}=\quad 5.8 \mathrm{Mpc} \quad 12+\log (0 / H)=7.74$
Morph Type: Sm-pec, BCDG, WR, HIIgalaxy
Other names : Haro29,


RA: 23h31m58.3s DEC: 28d56m50s
$D=77.8 \mathrm{Mpc} \quad 12+\log (0 / H)=8.03$
Morph Type : HIIgalaxy, merger,
Other names: N/A





NGC 1705


NGC 2366


| RA: 04h54m13.5s | DEC $:-53 \mathrm{~d} 21 \mathrm{~m} 40 \mathrm{~s}$ |
| :--- | :--- |
| $\mathrm{D}=$ | 5.1 Mpc |
| Morph Type : SA0-pec, HIIgalaxy, BCDG |  |
| Other names : ESO158, G013 |  |

NGC1705 in the Dwarf Galaxy Survey :

NGC2366 in the Dwarf Galaxy Survey :



| RA : 12h 15 m 39.2 s | DEC : 36d19m37s |
| :---: | :---: |
| $\mathrm{D}=2.9 \mathrm{Mpc}$ | $12+\log (\mathrm{O} / \mathrm{H})=8.26$ |
| Morph Type : IAB(s)m, HIIgalaxy <br> Other names : NGC4228, UGC07278 |  |
|  |  |




| RA : 12 h 28 m 11.1 s | DEC : 44d05m37s |
| :--- | :--- |
| D $=4.2 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=8.20$ |
| Morph Type : Ibm, HIIgalaxy, starburst |  |
| Other names : UGC07592 |  |
| NGC4449 in the Dwarf Galaxy Survey : |  |




| RA: 12h59m02.3s | DEC : 34d51m34s |
| :--- | :--- |
| D $=7.5 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=7.89$ |
| Morph Type : SB(s)m, HIIgalaxy, starburst |  |
| Other names : IC3961, IZw049, UGC08098 |  |


NGC 5253


RA: 01h35m04.6s DEC: -41d26m10s
$\mathrm{D}=3.9 \mathrm{Mpc} \quad 12+\log (\mathrm{O} / \mathrm{H})=8.22$

Morph Type : SB(s)m-sp, HIIgalaxy
Other names: ESO297, G005

NGC625 in the Dwarf Galaxy Survey :

$12+\log (0 / \mathrm{H})$

NGC6822 in the Dwarf Galaxy Survey :


Pox 186


| RA : 13 h 25 m 48.6 s | DEC $:-11 \mathrm{~d} 36 \mathrm{~m} 38 \mathrm{~s}$ |
| :--- | :--- |
| $\mathrm{D}=18.3 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=7.70$ |
| Morph Type : EmLineGalaxy |  |
| Other names : N/A |  |

Pox186 in the Dwarf Galaxy Survey :


SBS 0335-052


SBS0335-052 in the Dwarf Galaxy Survey :


| RA : 12h02m02.4s | DEC : 54d15m50s |
| :--- | :--- |
| D $=57.0 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=7.44$ |
| Morph Type : BCG |  |
| Other names : N/A |  |



SBS 1211+540



## SBS 1249+493




SBS 1415+437


SBS 1533+574


| RA : 15 h 34 m 13.8 s | DEC $: 57 \mathrm{~d} 17 \mathrm{~m} 06 \mathrm{~s}$ |
| :--- | :--- |
| $\mathrm{D}=54.2 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=8.05$ |
| Morph Type : BCG |  |
| Other names : NGC5965, viIZw611 |  |



## Tol 0618-402



| RA : 06h20m02.5s | DEC $:-40 \mathrm{~d} 18 \mathrm{m09s}$ |
| :--- | :--- |
| D $=150.8 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=8.09$ |
| Morph Type : BCG |  |
| Other names : N/A |  |

Tol0618-402 in the Dwarf Galaxy Survey :



## UGC 4483



R.A. (J2000)

| RA : 12 h 17 m 17.1 s | DEC $:-28 \mathrm{dO2m33s}$ |
| :--- | :---: |
| D $=120.5 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=7.52$ |
| Morph Type : BCG, HIIgalaxy |  |
| Other names : Tol00021 |  |



| RA : 08h37m03.0s | DEC : 69d46m31s |
| :--- | :--- |
| D $=3.2 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=7.46$ |
| Morph Type : BCG, HIIgalaxy |  |
| Other names : N/A |  |

UGC4483 in the Dwarf Galaxy Survey :



| RA: 01h43m14.7s | DEC : 19d58m32s |
| :--- | :--- |
| D $=11.0 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=7.50$ |
| Morph Type : Im |  |
| Other names : HS0140+1943 |  |




| RA: 01h44m41.3s | DEC : 4d53m26s |
| :--- | :--- |
| D $=22.7 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=7.82$ |
| Morph Type : HIIgalaxy |  |
| Other names : N/A |  |


UM 311


| RA: 01h 15 m 34.4 s | DEC : 0d51m46s |
| :--- | :--- |
| D $=23.5 \mathrm{Mpc}$ | $12+\log (0 / \mathrm{H})=8.36$ |
| Morph Type : HIIgalaxy, WR |  |
| Other names : N/A |  |




RA: 11h42m12.4s DEC: 0d20m03s
$D=87.8 \mathrm{Mpc} \quad 12+\log (0 / H)=8.32$ Morph Type : Sbpec, WR, HIIgalaxy, starburst

Other names : Mrk1304, UGC06665




| RA $: 11 \mathrm{~h} 51 \mathrm{~m} 33.3 \mathrm{~s}$ | DEC $:-2 \mathrm{~d} 22 \mathrm{~m} 22 \mathrm{~s}$ |
| :--- | :--- |
| $\mathrm{D}=13.2 \mathrm{Mpc}$ | $12+\log (\mathrm{O} / \mathrm{H})=7.73$ |
| Morph Type $: \mathrm{BCD}$ |  |
| Other names : N/A |  |

UM461 in the Dwarf Galaxy Survey :


VIIZw403 in the Dwarf Galaxy Survey :


## Appendix C

## IRS data for the DGS galaxies

Haro 11


Fig. C.1. See Fig. 6.6

Haro2


Haro3


He2-10


HSOO17 + 1055


HSOO52+2536


HS0822+3542


HS1222+3741



HS1304 + 3529


HS1319+3224


HS $1330+3651$



IZw18


IC10



IIZw40 ISOCAM


Mrk1089


Mrk1450
2 different aors for SL and LL


Mrk 153


Mrk209


Mrk930


NGC1140


## NGC1569



NGC1705


NGC2366


NGC4449


## NGC5253



Pox186


SBS0335-052


SBS1159+545


SBS1211+540


SBS1249+493


SBS1415+437
2 different aors for SL and LL



Tol0618-402


Tol1214-277


UGC4483


## UM 133



UM448



VIIZw403



Fig. C.2. See Fig. 6.5 (right)



















## Bibliography

Ackermann, M., Ajello, M., Baldini, L., et al. 2011, ApJ, 726, 81
Adamo, A., Östlin, G., Zackrisson, E., et al. 2010, MNRAS, 407, 870
Albrecht, M., Chini, R., Krügel, E., Müller, S. A. H., \& Lemke, R. 2004, A\&A, 414, 141
Albrecht, M., Krügel, E., \& Chini, R. 2007, A\&A, 462, 575
Aller, L. H., ed. 1984, Astrophysics and Space Science Library, Vol. 112, Physics of thermal gaseous nebulae
Aloisi, A., Clementini, G., Tosi, M., et al. 2007, ApJ, 667, L151
Aloisi, A., van der Marel, R. P., Mack, J., et al. 2005, ApJ, 631, L45
Amblard, A., Cooray, A., Serra, P., et al. 2010, A\&A, 518, L9
Anderson, L. D., Zavagno, A., Rodón, J. A., et al. 2010, A\&A, 518, L99
Aniano, G., Draine, B. T., Calzetti, D., et al. 2012, ApJ, 756, 138
Arendt, R. G., Odegard, N., Weiland, J. L., et al. 1998, ApJ, 508, 74
Asano, R. S., Takeuchi, T. T., Hirashita, H., \& Inoue, A. K. 2013, Earth, Planets, and Space, 65, 213
Asplund, M., Grevesse, N., Sauval, A. J., \& Scott, P. 2009, ARA\&A, 47, 481
Auld, R., Bianchi, S., Smith, M. W. L., et al. 2013, MNRAS, 428, 1880
Baas, F., Israel, F. P., \& Koornneef, J. 1994, A\&A, 284, 403
Bazell, D. \& Dwek, E. 1990, ApJ, 360, 142
Begum, A. \& Chengalur, J. N. 2005a, MNRAS, 362, 609
Begum, A. \& Chengalur, J. N. 2005b, MNRAS, 362, 609
Bendo, G. J., Dale, D. A., Draine, B. T., et al. 2006, ApJ, 652, 283
Bendo, G. J., Galliano, F., \& Madden, S. C. 2012, MNRAS, 423, 197
Bendo, G. J., Joseph, R. D., Wells, M., et al. 2003, AJ, 125, 2361
Bendo, G. J., Wilson, C. D., Pohlen, M., et al. 2010a, A\&A, 518, L65
Bendo, G. J., Wilson, C. D., Warren, B. E., et al. 2010b, MNRAS, 402, 1409
Bergvall, N., Zackrisson, E., Andersson, B.-G., et al. 2006, A\&A, 448, 513
Bernard, J.-P., Reach, W. T., Paradis, D., et al. 2008, AJ, 136, 919
Bettoni, D., Galletta, G., \& García-Burillo, S. 2003, A\&A, 405, 5

Bianchi, S. 2013, A\&A, 552, A89
Bocchio, M., Jones, A. P., Verstraete, L., et al. 2013, A\&A, 556, A6
Bocchio, M., Micelotta, E. R., Gautier, A.-L., \& Jones, A. P. 2012, A\&A, 545, A124
Bohren, C. F. \& Huffman, D. R. 1983, Absorption and scattering of light by small particles
Bolatto, A. D., Jackson, J. M., Wilson, C. D., \& Moriarty-Schieven, G. 2000, ApJ, 532, 909
Bolatto, A. D., Wolfire, M., \& Leroy, A. K. 2013, ARA\&A, 51, 207
Bonatto, C. J. \& Pastoriza, M. G. 1990, ApJ, 353, 445
Boquien, M., Calzetti, D., Combes, F., et al. 2011, AJ, 142, 111
Bordalo, V., Plana, H., \& Telles, E. 2009, ApJ, 696, 1668
Boselli, A., Ciesla, L., Buat, V., et al. 2010a, A\&A, 518, L61
Boselli, A., Ciesla, L., Cortese, L., et al. 2012, A\&A, 540, A54
Boselli, A., Eales, S., Cortese, L., et al. 2010b, PASP, 122, 261
Boselli, A., Lequeux, J., \& Gavazzi, G. 2002, A\&A, 384, 33
Bot, C., Ysard, N., Paradis, D., et al. 2010, A\&A, 523, A20+
Böttner, C., Klein, U., \& Heithausen, A. 2003, A\&A, 408, 493
Boulanger, F., Abergel, A., Bernard, J.-P., et al. 1996, A\&A, 312, 256
Braine, J., Duc, P.-A., Lisenfeld, U., et al. 2001, A\&A, 378, 51
Bresolin, F. 2007, ApJ, 656, 186
Brinks, E. \& Taylor, C. L. 1994, in European Southern Observatory Conference and Workshop Proceedings, Vol. 49, European Southern Observatory Conference and Workshop Proceedings, ed. G. Meylan \& P. Prugniel, 263

Brooks, A. M., Governato, F., Booth, C. M., et al. 2007, ApJ, 655, L17
Campbell, A., Terlevich, R., \& Melnick, J. 1986, MNRAS, 223, 811
Cannon, J. M., Dohm-Palmer, R. C., Skillman, E. D., et al. 2003, AJ, 126, 2806
Cannon, J. M., McClure-Griffiths, N. M., Skillman, E. D., \& Côté, S. 2004, ApJ, 607, 274
Cannon, J. M., Walter, F., Armus, L., et al. 2006, ApJ, 652, 1170
Cantalupo, C. M., Borrill, J. D., Jaffe, A. H., Kisner, T. S., \& Stompor, R. 2010, ApJS, 187, 212
Cardelli, J. A., Meyer, D. M., Jura, M., \& Savage, B. D. 1996, ApJ, 467, 334
Carignan, C. \& Beaulieu, S. 1989, ApJ, 347, 760
Carignan, C., Beaulieu, S., \& Freeman, K. C. 1990, AJ, 99, 178
Cen, R. \& Ostriker, J. P. 1999, ApJ, 519, L109
Chengalur, J. N., Pustilnik, S. A., Martin, J.-M., \& Kniazev, A. Y. 2006, MNRAS, 371, 1849
Ciesla, L., Boselli, A., Smith, M. W. L., et al. 2012, A\&A, 543, A161

Cohen, M., Wheaton, W. A., \& Megeath, S. T. 2003, AJ, 126, 1090
Combes, F. 1991, Galaxies et cosmologie
Condon, J. J. 1992, ARA\&A, 30, 575
Condon, J. J., Cotton, W. D., Greisen, E. W., et al. 1998, AJ, 115, 1693
Cormier, D., Lebouteiller, V., Madden, S. C., et al. 2012, A\&A, 548, A20
Cormier, D., Madden, S. C., Lebouteiller, V., et al. 2014, ArXiv 1401.0563
Cowie, L. L., Songaila, A., Hu, E. M., \& Cohen, J. G. 1996, AJ, 112, 839
Cresci, G., Mannucci, F., Sommariva, V., et al. 2012, MNRAS, 421, 262
Dale, D. A., Aniano, G., Engelbracht, C. W., et al. 2012, ApJ, 745, 95
Dale, D. A., Bendo, G. J., Engelbracht, C. W., et al. 2005, ApJ, 633, 857
Dale, D. A., Cohen, S. A., Johnson, L. C., et al. 2009, ApJ, 703, 517
Dale, D. A., Gil de Paz, A., Gordon, K. D., et al. 2007, ApJ, 655, 863
Dale, D. A. \& Helou, G. 2002, ApJ, 576, 159
Dale, D. A., Helou, G., Contursi, A., Silbermann, N. A., \& Kolhatkar, S. 2001a, ApJ, 549, 215
Dale, D. A., Helou, G., Neugebauer, G., et al. 2001b, AJ, 122, 1736
Davies, J. I., Bianchi, S., Cortese, L., et al. 2012, MNRAS, 419, 3505
Davoust, E. \& Contini, T. 2004, A\&A, 416, 515
de Blok, W. J. G. \& Walter, F. 2006, AJ, 131, 343
de Graauw, T., Helmich, F. P., Phillips, T. G., et al. 2010, A\&A, 518, L6
de Vaucouleurs, G., de Vaucouleurs, A., Corwin, Jr., H. G., et al. 1991, Third Reference Catalogue of Bright Galaxies, ed. de Vaucouleurs, G., de Vaucouleurs, A., Corwin, H. G., Jr., Buta, R. J., Paturel, G., \& Fouque, P. (de Vaucouleurs, G.)

Dekel, A. \& Silk, J. 1986, ApJ, 303, 39
Desert, F.-X., Boulanger, F., \& Puget, J. L. 1990, A\&A, 237, 215
Draine, B. T. 2003, ARA\&A, 41, 241
Draine, B. T., Dale, D. A., Bendo, G., et al. 2007, ApJ, 663, 866
Draine, B. T. \& Hensley, B. 2012, ApJ, 757, 103
Draine, B. T. \& Lazarian, A. 1998, ApJ, 508, 157
Draine, B. T. \& Lee, H. M. 1984, ApJ, 285, 89
Draine, B. T. \& Li, A. 2001, ApJ, 551, 807
Draine, B. T. \& Li, A. 2007, ApJ, 657, 810
Duc, P.-A. \& Mirabel, I. F. 1998, A\&A, 333, 813
Dudley, C. C. \& Wynn-Williams, C. G. 1997, ApJ, 488, 720

Duley, W. W. \& Williams, D. A. 1981, MNRAS, 196, 269
Dumke, M., Krause, M., \& Wielebinski, R. 2004, A\&A, 414, 475
Dunne, L., Gomez, H. L., da Cunha, E., et al. 2011, MNRAS, 417, 1510
Dupac, X., Bernard, J.-P., Boudet, N., et al. 2003, A\&A, 404, L11
Dwek, E. 1998, ApJ, 501, 643
Dwek, E., Arendt, R. G., Fixsen, D. J., et al. 1997, ApJ, 475, 565
Ehle, M. 2005, in The Magnetized Plasma in Galaxy Evolution, ed. K. T. Chyzy, K. Otmianowska-Mazur, M. Soida, \& R.-J. Dettmar, 31-38

Ekta, B. \& Chengalur, J. N. 2010, MNRAS, 403, 295
Ekta, B., Pustilnik, S. A., \& Chengalur, J. N. 2009, MNRAS, 397, 963
Engelbracht, C. W., Blaylock, M., Su, K. Y. L., et al. 2007, PASP, 119, 994
Engelbracht, C. W., Gordon, K. D., Rieke, G. H., et al. 2005, ApJ, 628, L29
Engelbracht, C. W., Rieke, G. H., Gordon, K. D., et al. 2008, ApJ, 678, 804
Erb, D. K., Shapley, A. E., Pettini, M., et al. 2006, ApJ, 644, 813
Eskew, M., Zaritsky, D., \& Meidt, S. 2012, AJ, 143, 139
Fathi, K., Axon, D. J., Storchi-Bergmann, T., et al. 2011, ApJ, 736, 77
Feldmann, R., Gnedin, N. Y., \& Kravtsov, A. V. 2012, ApJ, 747, 124
Ferrara, A. \& Dettmar, R.-J. 1994, ApJ, 427, 155
Fioc, M. \& Rocca-Volmerange, B. 1997, A\&A, 326, 950
Fitzpatrick, E. L. 1999, PASP, 111, 63
Fricke, K. J., Izotov, Y. I., Papaderos, P., Guseva, N. G., \& Thuan, T. X. 2001, AJ, 121, 169
Frye, B., Broadhurst, T., \& Benítez, N. 2002, ApJ, 568, 558
Galametz, M., Hony, S., Galliano, F., et al. 2013, MNRAS, 431, 1596
Galametz, M., Kennicutt, R. C., Albrecht, M., et al. 2012, MNRAS, 425, 763
Galametz, M., Madden, S., Galliano, F., et al. 2009, A\&A, 508, 645
Galametz, M., Madden, S. C., Galliano, F., et al. 2011, A\&A, 532, A56
Galametz, M., Madden, S. C., Galliano, F., et al. 2010, A\&A, 518, L55
Galliano, F., Dwek, E., \& Chanial, P. 2008, ApJ, 672, 214
Galliano, F., Hony, S., Bernard, J.-P., et al. 2011, A\&A, 536, A88
Galliano, F., Madden, S. C., Jones, A. P., Wilson, C. D., \& Bernard, J.-P. 2005, A\&A, 434, 867
Galliano, F., Madden, S. C., Jones, A. P., et al. 2003, A\&A, 407, 159
Garnett, D. R., Kennicutt, Jr., R. C., \& Bresolin, F. 2004, ApJ, 607, L21
Genzel, R., Tacconi, L. J., Rigopoulou, D., Lutz, D., \& Tecza, M. 2001, ApJ, 563, 527

Gieren, W., Pietrzyński, G., Nalewajko, K., et al. 2006, ApJ, 647, 1056
Gil de Paz, A. \& Madore, B. F. 2005, ApJS, 156, 345
Gil de Paz, A., Madore, B. F., \& Pevunova, O. 2003, ApJS, 147, 29
Glover, S. C. O. \& Clark, P. C. 2012, MNRAS, 421, 9
Gordon, D. \& Gottesman, S. T. 1981, AJ, 86, 161
Gordon, K. D., Clayton, G. C., Misselt, K. A., Landolt, A. U., \& Wolff, M. J. 2003, ApJ, 594, 279
Gordon, K. D., Engelbracht, C. W., Fadda, D., et al. 2007, PASP, 119, 1019
Gordon, K. D., Rieke, G. H., Engelbracht, C. W., et al. 2005, PASP, 117, 503
Gratier, P., Braine, J., Rodriguez-Fernandez, N. J., et al. 2010, A\&A, 512, A68
Grebel, E. K. 2001, in Dwarf galaxies and their environment, ed. K. S. de Boer, R.-J. Dettmar, \& U. Klein, 45

Grebel, E. K., Gallagher, III, J. S., \& Harbeck, D. 2003, AJ, 125, 1926
Grenier, I. A., Casandjian, J.-M., \& Terrier, R. 2005, Science, 307, 1292
Greve, A., Becker, R., Johansson, L. E. B., \& McKeith, C. D. 1996, A\&A, 312, 391
Griffin, M. J., Abergel, A., Abreu, A., et al. 2010, A\&A, 518, L3
Grocholski, A. J., van der Marel, R. P., Aloisi, A., et al. 2012, AJ, 143, 117
Grossi, M., Hunt, L. K., Madden, S., et al. 2010, A\&A, 518, L52
Groves, B., Dopita, M. A., Sutherland, R. S., et al. 2008, ApJS, 176, 438
Guelin, M., Zylka, R., Mezger, P. G., Haslam, C. G. T., \& Kreysa, E. 1995, A\&A, 298, L29
Guelin, M., Zylka, R., Mezger, P. G., et al. 1993, A\&A, 279, L37
Guhathakurta, P. \& Draine, B. T. 1989, ApJ, 345, 230
Guseva, N. G., Izotov, Y. I., Fricke, K. J., \& Henkel, C. 2012, A\&A, 541, A115
Guseva, N. G., Izotov, Y. I., Papaderos, P., \& Fricke, K. J. 2007, A\&A, 464, 885
Guseva, N. G., Izotov, Y. I., \& Thuan, T. X. 2000, ApJ, 531, 776
Guseva, N. G., Papaderos, P., Izotov, Y. I., et al. 2003a, A\&A, 407, 91
Guseva, N. G., Papaderos, P., Izotov, Y. I., et al. 2003b, A\&A, 407, 105
Guseva, N. G., Papaderos, P., Izotov, Y. I., Noeske, K. G., \& Fricke, K. J. 2004, A\&A, 421, 519
Haser, S. M., Pauldrach, A. W. A., Lennon, D. J., et al. 1998, A\&A, 330, 285
Helou, G. 1986, ApJ, 311, L33
Hermelo, I., Lisenfeld, U., Relaño, M., et al. 2013, A\&A, 549, A70
Higdon, S. J. U., Devost, D., Higdon, J. L., et al. 2004, PASP, 116, 975
Hirashita, H. 1999, ApJ, 522, 220
Hirashita, H. 2013, MNRAS, 429, 3390

Hirashita, H. \& Kuo, T.-M. 2011, MNRAS, 416, 1340
Hirashita, H., Tajiri, Y. Y., \& Kamaya, H. 2002, A\&A, 388, 439
Hodge, P., Smith, T., Eskridge, P., MacGillivray, H., \& Beard, S. 1991, ApJ, 379, 621
Holland, W. S., Robson, E. I., Gear, W. K., et al. 1999, MNRAS, 303, 659
Holweger, H. 2001, in American Institute of Physics Conference Series, Vol. 598, Joint SOHO/ACE workshop
"Solar and Galactic Composition", ed. R. F. Wimmer-Schweingruber, 23-30
Hony, S., Galliano, F., Madden, S. M., \& SAGE Consortium. 2010, in Bulletin of the American Astronomical Society, Vol. 42, American Astronomical Society Meeting Abstracts 215, 459.14

Hopkins, A. M., Schulte-Ladbeck, R. E., \& Drozdovsky, I. O. 2002, AJ, 124, 862
Huchtmeier, W. K. 1979, A\&A, 75, 170
Huchtmeier, W. K., Krishna, G., \& Petrosian, A. 2005, A\&A, 434, 887
Huchtmeier, W. K. \& Richter, O.-G. 1988, A\&A, 203, 237
Huchtmeier, W. K., Seiradakis, J. H., \& Materne, J. 1981, A\&A, 102, 134
Hunt, L., Bianchi, S., \& Maiolino, R. 2005, A\&A, 434, 849
Hunt, L. K., Thuan, T. X., Sauvage, M., \& Izotov, Y. I. 2006, ApJ, 653, 222
Hunter, D. 1997, PASP, 109, 937
Hunter, D. A., Elmegreen, B. G., \& Ludka, B. C. 2010, AJ, 139, 447
Hunter, D. A., Gallagher, III, J. S., Rice, W. L., \& Gillett, F. C. 1989, ApJ, 336, 152
Hunter, D. A. \& Kaufman, M. 2007, AJ, 134, 721
Hunter, D. A. \& Sage, L. 1993, PASP, 105, 374
Hunter, D. A., van Woerden, H., \& Gallagher, J. S. 1999, AJ, 118, 2184
Hunter, D. A., Zahedy, F., Bowsher, E. C., et al. 2011, AJ, 142, 173
Inoue, A. K. 2011, Earth, Planets, and Space, 63, 1027
Israel, F. P. 1997, A\&A, 328, 471
Israel, F. P. 2005, A\&A, 438, 855
Israel, F. P., Baas, F., Rudy, R. J., Skillman, E. D., \& Woodward, C. E. 2003, A\&A, 397, 87
Israel, F. P. \& de Bruyn, A. G. 1988, A\&A, 198, 109
Israel, F. P., Maloney, P. R., Geis, N., et al. 1996, ApJ, 465, 738
Israel, F. P., Wall, W. F., Raban, D., et al. 2010, A\&A, 519, A67
Issa, M. R., MacLaren, I., \& Wolfendale, A. W. 1990, A\&A, 236, 237
Izotov, I. I., Guseva, N. G., Lipovetskii, V. A., Kniazev, A. I., \& Stepanian, J. A. 1991, A\&A, 247, 303
Izotov, Y. I., Chaffee, F. H., Foltz, C. B., et al. 1999, ApJ, 527, 757
Izotov, Y. I., Papaderos, P., Guseva, N. G., Fricke, K. J., \& Thuan, T. X. 2004, A\&A, 421, 539

Izotov, Y. I., Stasińska, G., Meynet, G., Guseva, N. G., \& Thuan, T. X. 2006, A\&A, 448, 955
Izotov, Y. I. \& Thuan, T. X. 1998, ApJ, 500, 188
Izotov, Y. I. \& Thuan, T. X. 1999, ApJ, 511, 639
Izotov, Y. I., Thuan, T. X., \& Guseva, N. G. 2012, A\&A, 546, A122
Izotov, Y. I., Thuan, T. X., \& Lipovetsky, V. A. 1994, ApJ, 435, 647
Izotov, Y. I., Thuan, T. X., \& Lipovetsky, V. A. 1997, ApJS, 108, 1
Izotov, Y. I., Thuan, T. X., \& Stasińska, G. 2007, ApJ, 662, 15
James, A., Dunne, L., Eales, S., \& Edmunds, M. G. 2002, MNRAS, 335, 753
James, P. A. 1994, MNRAS, 269, 176
Jarrett, T. H., Chester, T., Cutri, R., Schneider, S. E., \& Huchra, J. P. 2003, AJ, 125, 525
Jarrett, T. H., Cohen, M., Masci, F., et al. 2011, ApJ, 735, 112
Jones, A. \& Tielens, A. 1994, in The Cold Universe, ed. T. Montmerle, C. J. Lada, I. F. Mirabel, \& J. Tran Thanh Van, 35

Jones, A. P. 2004, in Astronomical Society of the Pacific Conference Series, Vol. 309, Astrophysics of Dust, ed. A. N. Witt, G. C. Clayton, \& B. T. Draine, 347

Jones, A. P. 2012, A\&A, 545, C3
Jones, A. P., Tielens, A. G. G. M., \& Hollenbach, D. J. 1996, ApJ, 469, 740
Jones, A. P., Tielens, A. G. G. M., Hollenbach, D. J., \& McKee, C. F. 1997, in American Institute of Physics Conference Series, Vol. 402, American Institute of Physics Conference Series, ed. T. J. Bernatowicz \& E. Zinner, 595-613

Juvela, M. \& Ysard, N. 2012a, A\&A, 541, A33
Juvela, M. \& Ysard, N. 2012b, A\&A, 539, A71
Karachentsev, I. D., Dolphin, A. E., Geisler, D., et al. 2002, A\&A, 383, 125
Karachentsev, I. D., Karachentseva, V. E., Huchtmeier, W. K., \& Makarov, D. I. 2004, AJ, 127, 2031
Karachentsev, I. D., Sharina, M. E., Dolphin, A. E., et al. 2003, A\&A, 398, 467
Karczewski, O. Ł., Barlow, M. J., Page, M. J., et al. 2013, MNRAS, 431, 2493
Kelly, B. C., Shetty, R., Stutz, A. M., et al. 2012, ApJ, 752, 55
Kemper, F., Vriend, W. J., \& Tielens, A. G. G. M. 2004, ApJ, 609, 826
Kennicutt, R. C., Calzetti, D., Aniano, G., et al. 2011, PASP, 123, 1347
Kennicutt, Jr., R. C. 1989, ApJ, 344, 685
Kennicutt, Jr., R. C. 1998, ApJ, 498, 541
Kennicutt, Jr., R. C., Armus, L., Bendo, G., et al. 2003a, PASP, 115, 928
Kennicutt, Jr., R. C., Bresolin, F., \& Garnett, D. R. 2003b, ApJ, 591, 801
Kennicutt, Jr., R. C., Hao, C.-N., Calzetti, D., et al. 2009, ApJ, 703, 1672

Kepley, A. A., Zweibel, E. G., Wilcots, E. M., Johnson, K. E., \& Robishaw, T. 2011, ApJ, 736, 139
Kewley, L. J., Dopita, M. A., Sutherland, R. S., Heisler, C. A., \& Trevena, J. 2001, ApJ, 556, 121
Kewley, L. J. \& Ellison, S. L. 2008, ApJ, 681, 1183
Kim, M., Kim, E., Hwang, N., et al. 2009, ApJ, 703, 816
Kirkpatrick, A., Calzetti, D., Galametz, M., et al. 2013, ApJ, 778, 51
Kiuchi, G., Ohta, K., Sawicki, M., \& Allen, M. 2004, AJ, 128, 2743
Klein, U., Graeve, R., \& Wielebinski, R. 1983, A\&A, 117, 332
Kobulnicky, H. A., Dickey, J. M., Sargent, A. I., Hogg, D. E., \& Conti, P. S. 1995, AJ, 110, 116
Kobulnicky, H. A. \& Johnson, K. E. 1999, ApJ, 527, 154
Kobulnicky, H. A., Kennicutt, Jr., R. C., \& Pizagno, J. L. 1999, ApJ, 514, 544
Kobulnicky, H. A. \& Kewley, L. J. 2004, ApJ, 617, 240
Kobulnicky, H. A. \& Skillman, E. D. 1997, ApJ, 489, 636
Kong, X., Cheng, F. Z., Weiss, A., \& Charlot, S. 2002, A\&A, 396, 503
Köppen, J., Weidner, C., \& Kroupa, P. 2007, MNRAS, 375, 673
Kormendy, J. \& Sanders, D. B. 1992, ApJ, 390, L53
Krumholz, M. R. \& Gnedin, N. Y. 2011, ApJ, 729, 36
Kunth, D., Lequeux, J., Sargent, W. L. W., \& Viallefond, F. 1994, A\&A, 282, 709
Kunth, D. \& Östlin, G. 2000, A\&A Rev., 10, 1
Kunth, D. \& Sargent, W. L. W. 1986, ApJ, 300, 496
Laor, A. \& Draine, B. T. 1993, ApJ, 402, 441
Lebouteiller, V., Barry, D. J., Spoon, H. W. W., et al. 2011, ApJS, 196, 8
Lebouteiller, V., Bernard-Salas, J., Sloan, G. C., \& Barry, D. J. 2010, PASP, 122, 231
Lee, H. \& Skillman, E. D. 2004a, ApJ, 614, 698
Lee, H. \& Skillman, E. D. 2004b, ApJ, 614, 698
Lee, H., Skillman, E. D., \& Venn, K. A. 2006, ApJ, 642, 813
Lee, J. C., Salzer, J. J., Impey, C., Thuan, T. X., \& Gronwall, C. 2002, AJ, 124, 3088
Lee, M. G., Freedman, W., Mateo, M., et al. 1993, AJ, 106, 1420
Leger, A. \& Puget, J. L. 1984, A\&A, 137, L5
Legrand, F., Kunth, D., Roy, J.-R., Mas-Hesse, J. M., \& Walsh, J. R. 2000, A\&A, 355, 891
Lelli, F., Verheijen, M., Fraternali, F., \& Sancisi, R. 2012, A\&A, 537, A72
Lequeux, J., Peimbert, M., Rayo, J. F., Serrano, A., \& Torres-Peimbert, S. 1979, A\&A, 80, 155
Leroy, A., Bolatto, A., Walter, F., \& Blitz, L. 2006, ApJ, 643, 825
Leroy, A., Bolatto, A. D., Simon, J. D., \& Blitz, L. 2005, ApJ, 625, 763

Leroy, A., Cannon, J., Walter, F., Bolatto, A., \& Weiss, A. 2007, ApJ, 663, 990
Leroy, A. K., Bolatto, A., Gordon, K., et al. 2011, ApJ, 737, 12
Leroy, A. K., Walter, F., Bigiel, F., et al. 2009, AJ, 137, 4670
Levesque, E. M., Kewley, L. J., \& Larson, K. L. 2010, AJ, 139, 712
Li, A. \& Draine, B. T. 2001, ApJ, 554, 778
Liang, Y. C., Hammer, F., Yin, S. Y., et al. 2007, A\&A, 473, 411
Lisenfeld, U. \& Ferrara, A. 1998, ApJ, 496, 145
Lisenfeld, U., Israel, F. P., Stil, J. M., \& Sievers, A. 2002, A\&A, 382, 860
López-Sánchez, Á. R. \& Esteban, C. 2010, A\&A, 517, A85
López-Sánchez, Á. R., Esteban, C., \& Rodríguez, M. 2004, ApJS, 153, 243
López-Sánchez, Á. R., Koribalski, B. S., van Eymeren, J., et al. 2012, MNRAS, 419, 1051
Luridiana, V., Peimbert, A., Peimbert, M., \& Cerviño, M. 2003, ApJ, 592, 846
Lynds, R., Tolstoy, E., O’Neil, Jr., E. J., \& Hunter, D. A. 1998, AJ, 116, 146
Madden, S. C. 2000, New Astronomy Reviews, 44, 249
Madden, S. C., Galliano, F., Jones, A. P., \& Sauvage, M. 2006, A\&A, 446, 877
Madden, S. C., Poglitsch, A., Geis, N., Stacey, G. J., \& Townes, C. H. 1997, ApJ, 483, 200
Madden, S. C., Rémy, A., Galliano, F., et al. 2012, in IAU Symposium, Vol. 284, IAU Symposium, ed. R. J. Tuffs \& C. C. Popescu, 141-148

Madden, S. C., Rémy-Ruyer, A., Galametz, M., et al. 2013, PASP, 125, 600
Maddox, S. J., Efstathiou, G., Sutherland, W. J., \& Loveday, J. 1990, MNRAS, 243, 692
Magdis, G. E., Daddi, E., Béthermin, M., et al. 2012, ApJ, 760, 6
Magnelli, B., Saintonge, A., Lutz, D., et al. 2012, A\&A, 548, A22
Magrini, L., Bianchi, S., Corbelli, E., et al. 2011, A\&A, 535, A13
Magrini, L. \& Gonçalves, D. R. 2009, MNRAS, 398, 280
Maier, C., Meisenheimer, K., \& Hippelein, H. 2004, A\&A, 418, 475
Maiolino, R., Nagao, T., Grazian, A., et al. 2008, A\&A, 488, 463
Mannucci, F., Cresci, G., Maiolino, R., et al. 2009, MNRAS, 398, 1915
Marble, A. R., Engelbracht, C. W., van Zee, L., et al. 2010, ApJ, 715, 506
Marzke, R. O. \& da Costa, L. N. 1997, AJ, 113, 185
Masegosa, J., Moles, M., \& Campos-Aguilar, A. 1994, ApJ, 420, 576
Mateo, M. L. 1998, ARA\&A, 36, 435
Mather, J. C., Cheng, E. S., Cottingham, D. A., et al. 1994, ApJ, 420, 439
Mathis, J. S. 1990, in Astronomical Society of the Pacific Conference Series, Vol. 12, The Evolution of the Interstellar Medium, ed. L. Blitz, 63-77

Mathis, J. S., Mezger, P. G., \& Panagia, N. 1983, A\&A, 128, 212
Mathis, J. S., Rumpl, W., \& Nordsieck, K. H. 1977, ApJ, 217, 425
Mauch, T., Murphy, T., Buttery, H. J., et al. 2008, VizieR Online Data Catalog, 8081, 0
McCall, M. L. 1984, MNRAS, 208, 253
McCall, M. L., Rybski, P. M., \& Shields, G. A. 1985, ApJS, 57, 1
McConnachie, A. W. 2012, AJ, 144, 4
McGaugh, S. S. 1991, ApJ, 380, 140
McQuade, K., Calzetti, D., \& Kinney, A. L. 1995, ApJS, 97, 331
Meier, D. S., Turner, J. L., Crosthwaite, L. P., \& Beck, S. C. 2001, AJ, 121, 740
Meixner, M., Galliano, F., Hony, S., et al. 2010, A\&A, 518, L71
Melisse, J. P. M. \& Israel, F. P. 1994, A\&A, 285, 51
Meny, C., Gromov, V., Boudet, N., et al. 2007, A\&A, 468, 171
Meurer, G. R., Staveley-Smith, L., \& Killeen, N. E. B. 1998, MNRAS, 300, 705
Meyer, D. M., Cardelli, J. A., \& Sofia, U. J. 1997, ApJ, 490, L103
Meyer, D. M., Jura, M., \& Cardelli, J. A. 1998, ApJ, 493, 222
Moles, M., Marquez, I., Masegosa, J., et al. 1994, ApJ, 432, 135
Moll, S. L., Mengel, S., de Grijs, R., Smith, L. J., \& Crowther, P. A. 2007, MNRAS, 382, 1877
Moore, B., Ghigna, S., Governato, F., et al. 1999, ApJ, 524, L19
Mould, J. R., Huchra, J. P., Freedman, W. L., et al. 2000, ApJ, 529, 786
Moustakas, J., Kennicutt, Jr., R. C., Tremonti, C. A., et al. 2010, ApJS, 190, 233
Muñoz-Mateos, J. C., Gil de Paz, A., Boissier, S., et al. 2009, ApJ, 701, 1965
Murphy, E. J., Helou, G., Condon, J. J., et al. 2010, ApJ, 709, L108
Navarro, J. F., Frenk, C. S., \& White, S. D. M. 1996, ApJ, 462, 563
Neininger, N., Guelin, M., Garcia-Burillo, S., Zylka, R., \& Wielebinski, R. 1996, A\&A, 310, 725
Oh, S.-H., de Blok, W. J. G., Brinks, E., Walter, F., \& Kennicutt, Jr., R. C. 2011, AJ, 141, 193
O'Halloran, B., Galametz, M., Madden, S. C., et al. 2010, A\&A, 518, L58
O'Halloran, B., Satyapal, S., \& Dudik, R. P. 2006, ApJ, 641, 795
Olive, K. A., Steigman, G., \& Skillman, E. D. 1997, ApJ, 483, 788
Ott, S. 2010, in Astronomical Society of the Pacific Conference Series, Vol. 434, Astronomical Data Analysis Software and Systems XIX, ed. Y. Mizumoto, K.-I. Morita, \& M. Ohishi, 139

Pagel, B. E. J. 1997, Nucleosynthesis and Chemical Evolution of Galaxies
Pagel, B. E. J., Simonson, E. A., Terlevich, R. J., \& Edmunds, M. G. 1992, MNRAS, 255, 325
Paradis, D., Bernard, J.-P., \& Mény, C. 2009, A\&A, 506, 745

Paradis, D., Veneziani, M., Noriega-Crespo, A., et al. 2010, A\&A, 520, L8
Parkin, T. J., Wilson, C. D., Foyle, K., et al. 2012, MNRAS, 422, 2291
Paturel, G., Theureau, G., Bottinelli, L., et al. 2003, A\&A, 412, 57
Peeters, E., Hony, S., Van Kerckhoven, C., et al. 2002, A\&A, 390, 1089
Peimbert, A., Peimbert, M., \& Luridiana, V. 2002, ApJ, 565, 668
Persic, M., Mariani, S., Cappi, M., et al. 1998, A\&A, 339, L33
Petitpas, G. R. \& Wilson, C. D. 2003, ApJ, 587, 649
Pettini, M. \& Pagel, B. E. J. 2004, MNRAS, 348, L59
Pilbratt, G. L., Riedinger, J. R., Passvogel, T., et al. 2010, A\&A, 518, L1
Pilleri, P., Montillaud, J., Berné, O., \& Joblin, C. 2012, A\&A, 542, A69
Pilyugin, L. S. 2001, A\&A, 374, 412
Pilyugin, L. S. \& Thuan, T. X. 2005, ApJ, 631, 231
Pilyugin, L. S. \& Thuan, T. X. 2007, ApJ, 669, 299
Pilyugin, L. S., Thuan, T. X., \& Vílchez, J. M. 2006, MNRAS, 367, 1139
Planck Collaboration, Ade, P. A. R., Aghanim, N., et al. 2011a, A\&A, 536, A19
Planck Collaboration, Ade, P. A. R., Aghanim, N., et al. 2011b, A\&A, 536, A20
Poglitsch, A., Krabbe, A., Madden, S. C., et al. 1995, ApJ, 454, 293
Poglitsch, A., Waelkens, C., Geis, N., et al. 2010, A\&A, 518, L2
Popescu, C. C. \& Hopp, U. 2000, A\&AS, 142, 247
Pustilnik, S. A., Kniazev, A. Y., Pramskij, A. G., Ugryumov, A. V., \& Masegosa, J. 2003, A\&A, 409, 917
Pustilnik, S. A. \& Martin, J.-M. 2007, A\&A, 464, 859
Pustilnik, S. A., Martin, J.-M., Huchtmeier, W. K., et al. 2002, A\&A, 389, 405
Reach, W. T., Dwek, E., Fixsen, D. J., et al. 1995, ApJ, 451, 188
Reach, W. T., Megeath, S. T., Cohen, M., et al. 2005, PASP, 117, 978
Rémy-Ruyer, A., Madden, S. C., Galliano, F., et al. 2013, A\&A, 557, A95
Revaz, Y. \& Jablonka, P. 2012, A\&A, 538, A82
Rice, W., Lonsdale, C. J., Soifer, B. T., et al. 1988, ApJS, 68, 91
Rieke, G., Lada, C., Lebofsky, M., et al. 1984, in Bulletin of the American Astronomical Society, Vol. 16, Bulletin of the American Astronomical Society, 906

Roman-Duval, J., Israel, F. P., Bolatto, A., et al. 2010, A\&A, 518, L74+
Rosenberg, J. L., Ashby, M. L. N., Salzer, J. J., \& Huang, J.-S. 2006, ApJ, 636, 742
Roussel, H. 2013, PASP, 125, 1126
Sage, L. J., Salzer, J. J., Loose, H.-H., \& Henkel, C. 1992, A\&A, 265, 19

Sage, L. J., Welch, G. A., \& Young, L. M. 2007, ApJ, 657, 232
Sanders, D. B., Mazzarella, J. M., Kim, D.-C., Surace, J. A., \& Soifer, B. T. 2003, AJ, 126, 1607
Sandstrom, K. M., Leroy, A. K., Walter, F., et al. 2013, ApJ, 777, 5
Sauvage, M., Thuan, T. X., \& Lagage, P. O. 1997, A\&A, 325, 98
Sauvage, M., Vigroux, L., \& Thuan, T. X. 1990, A\&A, 237, 296
Savage, B. D. \& Sembach, K. R. 1996, ARA\&A, 34, 279
Savaglio, S., Glazebrook, K., Le Borgne, D., et al. 2005, ApJ, 635, 260
Saviane, I., Ivanov, V. D., Held, E. V., et al. 2008, A\&A, 487, 901
Schruba, A., Leroy, A. K., Walter, F., et al. 2012, AJ, 143, 138
Schuller, F. 2012, in Society of Photo-Optical Instrumentation Engineers (SPIE) Conference Series, Vol. 8452, Society of Photo-Optical Instrumentation Engineers (SPIE) Conference Series

Schulte-Ladbeck, R. E., Hopp, U., Greggio, L., Crone, M. M., \& Drozdovsky, I. O. 2001, AJ, 121, 3007
Searle, L. \& Sargent, W. L. W. 1972, ApJ, 173, 25
Serra Díaz-Cano, L. \& Jones, A. P. 2008, A\&A, 492, 127
Sharina, M. E., Karachentsev, I. D., \& Tikhonov, N. A. 1996, A\&AS, 119, 499
Shetty, R., Glover, S. C., Dullemond, C. P., et al. 2011, MNRAS, 415, 3253
Shetty, R., Kauffmann, J., Schnee, S., \& Goodman, A. A. 2009a, ApJ, 696, 676
Shetty, R., Kauffmann, J., Schnee, S., Goodman, A. A., \& Ercolano, B. 2009b, ApJ, 696, 2234
Simon, J. D. \& Geha, M. 2007, ApJ, 670, 313
Simpson, C. E. \& Gottesman, S. T. 2000, AJ, 120, 2975
Skibba, R. A., Engelbracht, C. W., Dale, D., et al. 2011, ApJ, 738, 89
Skillman, E. D., Côté, S., \& Miller, B. W. 2003, AJ, 125, 610
Smith, J. D. T., Draine, B. T., Dale, D. A., et al. 2007, ApJ, 656, 770
Smith, M. W. L., Eales, S. A., Gomez, H. L., et al. 2012a, ApJ, 756, 40
Smith, M. W. L., Gomez, H. L., Eales, S. A., et al. 2012b, ApJ, 748, 123
Smith, R. K. \& Dwek, E. 1998, ApJ, 503, 831
Smoker, J. V., Davies, R. D., Axon, D. J., \& Hummel, E. 2000, A\&A, 361, 19
Snow, T. P. \& Witt, A. N. 1996, ApJ, 468, L65
Sofia, U. J. \& Meyer, D. M. 2001, ApJ, 554, L221
Sramek, R. A. \& Weedman, D. W. 1986, ApJ, 302, 640
Stansberry, J. A., Gordon, K. D., Bhattacharya, B., et al. 2007, PASP, 119, 1038
Stasińska, G. 1990, A\&AS, 83, 501
Stasińska, G. 2005, A\&A, 434, 507

Stasińska, G. \& Izotov, Y. 2003, A\&A, 397, 71
Steigman, G. 2005, Physica Scripta Volume T, 121, 142
Stepnik, B., Abergel, A., Bernard, J., et al. 2003, A\&A, 398, 551
Stepnik, B., Abergel, A., Bernard, J.-P., et al. 2001, in Astronomical Society of the Pacific Conference Series, Vol. 243, From Darkness to Light: Origin and Evolution of Young Stellar Clusters, ed. T. Montmerle \& P. André, 47

Strong, A. W., Moskalenko, I. V., Reimer, O., Digel, S., \& Diehl, R. 2004, A\&A, 422, L47
Swaters, R. A., van Albada, T. S., van der Hulst, J. M., \& Sancisi, R. 2002, A\&A, 390, 829
Taylor, C. L., Kobulnicky, H. A., \& Skillman, E. D. 1998, AJ, 116, 2746
Thuan, T. X. 1983, ApJ, 268, 667
Thuan, T. X. 1985, ApJ, 299, 881
Thuan, T. X., Hibbard, J. E., \& Lévrier, F. 2004, AJ, 128, 617
Thuan, T. X., Izotov, Y. I., \& Lipovetsky, V. A. 1995, ApJ, 445, 108
Thuan, T. X., Lipovetsky, V. A., Martin, J.-M., \& Pustilnik, S. A. 1999a, A\&AS, 139, 1
Thuan, T. X. \& Martin, G. E. 1981, ApJ, 247, 823
Thuan, T. X., Sauvage, M., \& Madden, S. 1999b, ApJ, 516, 783
Tielens, A. G. G. M. 2005, The Physics and Chemistry of the Interstellar Medium (Cambridge University Press)

Tielens, A. G. G. M., McKee, C. F., Seab, C. G., \& Hollenbach, D. J. 1994, ApJ, 431, 321
Tolstoy, E., Hill, V., \& Tosi, M. 2009, ARA\&A, 47, 371
Tosi, M., Sabbi, E., Bellazzini, M., et al. 2001, AJ, 122, 1271
Tremonti, C. A., Heckman, T. M., Kauffmann, G., et al. 2004, ApJ, 613, 898
Tully, R. B. 1988, Nearby galaxies catalog (Cambridge University Press)
Tully, R. B., Bottinelli, L., Fisher, J. R., et al. 1978, A\&A, 63, 37
Ugryumov, A. V., Engels, D., Kniazev, A. Y., et al. 2001, A\&A, 374, 907
Ugryumov, A. V., Engels, D., Pustilnik, S. A., et al. 2003, A\&A, 397, 463
Urry, C. M. \& Padovani, P. 1995, PASP, 107, 803
Valcke, S., de Rijcke, S., \& Dejonghe, H. 2008, MNRAS, 389, 1111
van den Bergh, S. 2009, ApJ, 694, L120
van Eymeren, J., Marcelin, M., Koribalski, B. S., et al. 2009, A\&A, 505, 105
van Zee, L. \& Haynes, M. P. 2006, ApJ, 636, 214
van Zee, L., Haynes, M. P., Salzer, J. J., \& Broeils, A. H. 1996, AJ, 112, 129
van Zee, L., Salzer, J. J., \& Skillman, E. D. 2001, AJ, 122, 121
van Zee, L., Westpfahl, D., Haynes, M. P., \& Salzer, J. J. 1998, AJ, 115, 1000

Veilleux, S., Kim, D.-C., Sanders, D. B., Mazzarella, J. M., \& Soifer, B. T. 1995, ApJS, 98, 171
Vennik, J., Hopp, U., \& Popescu, C. C. 2000, A\&AS, 142, 399
Walter, F., Brinks, E., de Blok, W. J. G., et al. 2008, AJ, 136, 2563
Walter, F., Cannon, J. M., Roussel, H., et al. 2007, ApJ, 661, 102
Walter, F., Taylor, C. L., Hüttemeister, S., Scoville, N., \& McIntyre, V. 2001, AJ, 121, 727
Weingartner, J. C. \& Draine, B. T. 2001, ApJ, 548, 296
Welch, G. A. \& Sage, L. J. 2003, ApJ, 584, 260
Welty, D. E., Jenkins, E. B., Raymond, J. C., Mallouris, C., \& York, D. G. 2002, ApJ, 579, 304
Whittet, D. C. B. 2003, Astronomy and Geophysics, 44, 35
Williams, B. A., McMahon, P. M., \& van Gorkom, J. H. 1991, AJ, 101, 1957
Wilson, C. D. 1995, ApJ, 448, L97
Wilson, C. D., Warren, B. E., Israel, F. P., et al. 2012, MNRAS, 424, 3050
Wolfire, M. G., Hollenbach, D., \& McKee, C. F. 2010, ApJ, 716, 1191
Worthey, G. 1994, ApJS, 95, 107
Wright, E. L., Eisenhardt, P. R. M., Mainzer, A. K., et al. 2010, AJ, 140, 1868
Wu, Y., Charmandaris, V., Hao, L., et al. 2006, ApJ, 639, 157
Yang, M. \& Phillips, T. 2007, ApJ, 662, 284
Yin, S. Y., Liang, Y. C., \& Zhang, B. 2007, in Astronomical Society of the Pacific Conference Series, Vol. 373, The Central Engine of Active Galactic Nuclei, ed. L. C. Ho \& J.-W. Wang, 686

Young, J. S. \& Knezek, P. M. 1989, ApJ, 347, L55
Young, J. S., Xie, S., Tacconi, L., et al. 1995, ApJS, 98, 219
Young, L. M., Bureau, M., Davis, T. A., et al. 2011, MNRAS, 414, 940
Young, L. M. \& Lo, K. Y. 1997, ApJ, 490, 710
Ysard, N., Juvela, M., Demyk, K., et al. 2012, A\&A, 542, A21
Ysard, N. \& Verstraete, L. 2010, A\&A, 509, A12
Zhao, Y., Gao, Y., \& Gu, Q. 2013, ApJ, 764, 44
Zhu, M., Papadopoulos, P. P., Xilouris, E. M., Kuno, N., \& Lisenfeld, U. 2009, ApJ, 706, 941
Zhukovska, S. 2008, PhD thesis, Ruperto-Carola-University
Zhukovska, S. 2014, ArXiv 1401.1675
Zhukovska, S., Gail, H.-P., \& Trieloff, M. 2008, A\&A, 479, 453
Zubko, V., Dwek, E., \& Arendt, R. G. 2004, ApJS, 152, 211


[^0]:    ${ }^{1}$ Images from http://www.sedfitting.org/SED08/Welcome.html

[^1]:    ${ }^{2}$ Image from http://www.astro.ucla.edu/ ghezgroup/gc/journey/wave.shtml

[^2]:    ${ }^{3}$ The albedo of a dust grain describes the relative importance of scattering compared to absorption by the dust grain.

[^3]:    ${ }^{4}$ The energy density is related to the mean intensity of a given radiation by: $u_{\lambda}=\frac{4 \pi}{c} \times J_{\lambda}$

[^4]:    ${ }^{5} 1$ Habing $=1.6 \times 10^{-3} \mathrm{erg} \mathrm{cm}^{-2} \mathrm{~s}^{-1}$

[^5]:    ${ }^{6}$ Note that this is not the same galaxy, nor the same model as in Fig. 1.2.

[^6]:    ${ }^{1} R_{25}$ is defined as the radius at which the surface brightness falls to a level of $25 \mathrm{mag} / \mathrm{arcsec}{ }^{2}$.

[^7]:    ${ }^{2}$ Image from http://cseligman.com/text/atlas/
    ${ }^{3}$ Image from http://billsnyderastrophotography.com

[^8]:    ${ }^{4}$ Image from http://hubblesite.org

[^9]:    ${ }^{5}$ Images of the dSph and dE from http://www.noao.edu/outreach/aop/observers/
    ${ }^{6}$ Image from Braine et al. (2001)

[^10]:    ${ }^{1}$ Image from the Herschel Observers' Manual, available at http://herschel.esac.esa.int/twiki/bin/view/Public/ SpacecraftObservatoryWeb?template=viewprint.

[^11]:    ${ }^{2}$ Image from the PACS Observers' Manual, available at http://herschel.esac.esa.int/twiki/bin/view/Public/ PacsCalibrationWeb?template=viewprint.

[^12]:    ${ }^{3}$ Image from the PACS Observers' Manual.

[^13]:    ${ }^{4}$ Image from the SPIRE Observers' Manual, available at http://herschel.esac.esa.int/twiki/bin/view/Public/ SpireCalibrationWeb?template=viewprint.
    ${ }^{5}$ Since this manuscript has been written, the values for the SPIRE calibration and the uncertainties may have been updated. The same goes for the beam areas. We advise the reader to treat these numbers with caution.

[^14]:    ${ }^{6}$ Image from the SPIRE Observers' Manual.

[^15]:    ${ }^{7}$ Remember from Chapter 2, $R_{23}=\frac{[O I I] \lambda 3727+[O I I I] \lambda \lambda 4959,5007}{H \beta}$

[^16]:    ${ }^{1}$ However, the reader must keep in mind that the current version of HIPE might not be the same than that used for the data reduction and photometry in this Chapter and that the maps and flux densities might need some updating. The flux densities presented here are the same as in Rémy-Ruyer et al. (2013).
    ${ }^{2}$ The version 7 cited here corresponds to the value of the CALFileVersion metadata of the Responsitivity Calibration Product in HIPE.

[^17]:    ${ }^{3}$ The version 5 of HIPE was used for producing the SPIRE maps.

[^18]:    ${ }^{4}$ The last version of this source fitter is incorporated into HIPE v10 (Bendo et al. in prep.).

[^19]:    ${ }^{5}$ SPIRE photometre reference spectrum values: http://herschel.esac.esa.int/twiki/bin/view/Public/SpirePhotometer BeamProfileAnalysis, September 2012 values.

[^20]:    ${ }^{6}$ This value is given in: http://herschel.esac.esa.int/twiki/bin/view/Public/SpirePhotometerBeamProfileAnalysis.

[^21]:    ${ }^{1}$ for their "Standard Model", see Appendix A of Galliano et al. (2011).

[^22]:    ${ }^{2}$ This documentation is available at: http://herschel.esac.esa.int/twiki/bin/view/Public/PacsCalibrationWeb? template=viewprint
    ${ }^{3}$ This value is given in: http://herschel.esac.esa.int/twiki/bin/view/Public/SpirePhotometerBeamProfileAnalysis.

[^23]:    ${ }^{4}$ The Spearman rank coefficient, $\rho$, indicates how well the relationship between X and Y can be described by a monotonic function: monotonically increasing: $\rho>0$, or monotonically decreasing: $\rho<0$.

[^24]:    ${ }^{1}$ The query form is available at: http://sha.ipac.caltech.edu/applications/Spitzer/SHA/\#id=SearchByPosition\& startIdx $=0$ \&pageSize $=0$ \&shortDesc $=$ Position\%20Search\&isBookmarkAble=true

[^25]:    ${ }^{a}$ : Extended aperture corrections coefficients, see Section 6.2.1, Photometry.

[^26]:    ${ }^{2}$ This document is available at http://irsa.ipac.caltech.edu/data/SPITZER/docs/irac/iracinstrumenthandbook/

[^27]:    ${ }^{3}$ see the IRS Handbook, available at http://irsa.ipac.caltech.edu/data/SPITZER/docs/irs/irsinstrumenthandbook/

[^28]:    ${ }^{4}$ http://cassis.astro.cornell.edu/atlas/

[^29]:    ${ }^{5}$ The corresponding documentation for PACS colour corrections is available at http://herschel.esac.esa.int/twiki/ pub/Public/PacsCalibrationWeb/cc_report_v1.pdf. The values presented in this documentation can also be rederived from Eq. 3.8.
    ${ }^{6}$ The MIPS Instrument Handbook is available at http://irsa.ipac.caltech.edu/data/SPITZER/docs/mips/ mipsinstrumenthandbook/home/

[^30]:    ${ }^{7}$ available at: http://irsa.ipac.caltech.edu/cgi-bin/Gator/nph-dd
    ${ }^{8}$ available at: http://wise2.ipac.caltech.edu/docs/release/allsky/expsup/index.html

[^31]:    ${ }^{9}$ http://www.astro.uni-bonn.de/boawiki/Boa

[^32]:    $a$ : The aperture is the same as the one used for Herschel.
    ${ }^{b}$ : The Herschel aperture has been shorten to avoid a contaminating source.
    ${ }^{c}$ : The Herschel aperture has been adapted to match the peculiar morphology of the source in the NIR. The radius indicated in between parenthesis is the equivalent radius $\mathrm{r}=\sqrt{a b}$
    to be used in Eq. 6.10 .
    $e$ : The IRAC map is too small for the Herschel aperture. The aperture thus had to be shorten.
    $f$ : For this galaxy, the IRAC maps do not cover the full galaxy. Thus we do not feel comfortable in reporting flux densities for this source.
    $g$ : This flux density is more coherent with the rest of the IR photometry for this galaxy (see Section 7.2.4). Thus we recommend to use this value when studying this object.

[^33]:    This source was not found in the database.
    No good measurement was available from the database as the source is too extended or present very low surface brightness areas. ${ }^{c}$ : This source is extended but no $\mathrm{W}_{i}$ MAG measurement was available.
    $d$ : The photometry is not done on the whole complex and thus we do not report it here.
    : The aperture used by the database do not encompass the total galaxy at these wavelengths.

[^34]:    ${ }^{1}$ We adopt the notation $F_{i}$ instead of $f_{i}$ in Zubko et al. (2004) to avoid confusion with the parameters $f_{P A H}$ and $f_{s i l}$ of the model, see Section 7.2.1.
    ${ }^{2}$ We adopt the notation $S_{\lambda}^{i}(\lambda, a)$ instead of $E_{\lambda}^{i}(\lambda, a)$ as in Zubko et al. (2004) because this quantity actually correspond to the source function: $S_{\lambda}=j_{\lambda} / \alpha_{\lambda}$ rather than to an "emissivity" as it is called in Zubko et al. (2004). For isotropic emission, the emissivity is defined as: $\epsilon_{\lambda}=\frac{4 \pi j_{\lambda}(\lambda)}{\rho_{I S M}}$.

[^35]:    ${ }^{3}$ COsmic Background Explorer.
    ${ }^{4}$ DIRBE (Diffuse InfraRed Background Experiment) and FIRAS (Far InfraRed Absolute Spectrometer) are two instruments on board COBE.

[^36]:    ${ }^{a}$ : This reference is used by Zubko et al. (2004) in his model to get the size distribution.
    ${ }^{b}$ : This reference is the updated version by Draine \& Li (2007) of the PAH optical properties that are used in our model to compute the grain absorption cross-sections.

[^37]:    ${ }^{5}$ Notations: $L_{\nu}$ is the monochromatic luminosity and $I_{\lambda}$ from Eq. 7.3 is the specific monochromatic intensity per unit solid angle: $L_{\nu}=4 \pi I_{\nu}=4 \pi \frac{\lambda}{\nu} I_{\lambda}$. The monochromatic luminosities per unit dust mass of a given dust specie, $l_{\nu}^{i}$, is given by $l_{\nu}^{i}=L_{\nu}^{i} / \mathrm{M}_{i}$.

[^38]:    ${ }^{6}$ The enthalpy is linked to the heat capacity, $\mathrm{C}(\mathrm{T})$, defined in Section 1.2 .4 by: $\mathrm{H}(\mathrm{T})=\frac{4}{3} \pi a^{3} \int_{0}^{T} C\left(T_{g}\right) \mathrm{d} T_{g}$ where $T_{g}$ is the grain temperature.

[^39]:    ${ }^{7}$ Projet d'Étude des GAlaxies par Synthèse Évolutive

[^40]:    ${ }^{8}$ Available at: http://lambda.gsfc.nasa.gov/product/iras/docs/exp.sup/

[^41]:    ${ }^{9}$ Note the very large error bar for the dust mass due to the large error bar on the SPIRE observations.

[^42]:    *: 0.00 for the minimum $f_{P A H}$ value means that no PAHs are detected in at least one galaxy in the sample.

[^43]:    10 100
    Wavelength [ $\mu \mathrm{m}$ ]

[^44]:    ${ }^{1}$ as of January2014.
    ${ }^{2} 2014$ or 2015 .

[^45]:    ${ }^{3}$ The TIR luminosity is defined in Eq. 7.26 and estimated between 1 and $1000 \mu \mathrm{~m}$.

[^46]:    ${ }^{4}$ We remind the reader here that non-dust contaminations have already been removed from the total submm flux densities at 850,870 and $1200 \mu \mathrm{~m}$ (see Section 6.5).

[^47]:    ${ }^{1}$ This value is from Table 6 from Zubko et al. (2004) for the BARE-GR-S model, which corresponds to the dust composition used for our modelling (see Section 7.1).
    ${ }^{2}$ Remember: $\mathrm{R}_{23}=([\mathrm{OII}] \lambda 3727+[\mathrm{OIII}] \lambda \lambda 4959,5007) / \mathrm{H} \beta$.

[^48]:    ${ }^{3}$ http://www.physics.wisc.edu/ craigm/idl/idl.html.

[^49]:    ${ }^{4}$ This updated SPIRE calibration from September 2012 had the effect of decreasing the SPIRE flux densities by about $10 \%$ compared to the Galliano et al. (2011) study.

[^50]:    ${ }^{5}$ The star-formation timescale, $\tau_{\mathrm{SF}}$, is defined by the timescale during which star formation occurs: $\tau_{\mathrm{SF}}=\left(M_{\mathrm{ISM}}\right) / \mathrm{SFR}$, where SFR is the star-formation rate (see Eq. 5 of Asano et al. 2013).

