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Fusion des galaxies juvéniles : des simulations aux observations

Valentin Perret

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Valentin PERRET

Fusions de galaxies juvéniles Des simulations aux observations

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Table des acronymes

AGN	Active Galaxy Nuclei
ALMA	Atacama Large Millimeter/submillimeter Array
AMR	Adaptive Mesh Refinement
CCD	Charge-Coupled Device
CFHT	Canada-France-Hawaiï Telescope
CPU	Central Processing Unit
DM	Dark Matter
DICE	Disk Initial Conditions Environment
ELT	Extremely Large Telescope
ESO	European Southern Observatory
FFT	Fast Fourier Transform
FWHM	Full Width at Half Maximum
GADGET	GAxaxies with Dark matter and Gas intEracT
GALMER	GALaxy MERgers database
GB	Giga Byte
IDL	Interactive Data Languge
IFU	Integral Field Units
IC	Initial Conditions
IGM	Inter-Galactic Medium
ISM	Inter-Stellar Medium
IMAGES	Intermediate MAAss Galaxy Evolution Sequence
KINEGEN	KINEmetry in GENetic algorithm
KMOS	K-band Multi-Object Spectrograph
LASER	Light Amplification by the Stimulated Emission of Radiation
LIRG	Luminous Infra-Red Galaxie
MASSIV	Mass ASsembly with SInfoni and Vvds
MCMC	Monte-Carlo Markov Chain
MIRAGE	Merging and Isolated high-Redshit Amr GalaxiEs
MOND	Modified Newtonian Dynamics
MOSFIRE	Multi-Object Spectrometer for Infra-Red Exploration
NOEMA	Northern Extended Millimeter Array
AO	Adaptive Optics
OSIRIS	OH-Suppressing Infra-Red Imaging Spectrograph
PA	Position Angle of the major axis
PSF	Point Spread Function en anglais, Réponse Impulsionnelle en français
RAM	Random Access Memory
SFR	Star Formation Rate
SINFONI	Spectrograph for INtegral Field Observations in the Near Infrared

SINS	Spectroscopic Imaging survey in the Near-infrared
SKA	Square Kilometer Array
SPH	Smoothed Particles Hydrodynamics
TFR	Tully-Fischer Relation
VIMOS	Visible MultiObject Spectrograph
VLT	Very Large Telescope
VVDS	Vimos-Vlt Deep Survey
WMAP	Wilkinson Microwave Anisotropy Probe

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Prélude

La Voie Lactée

Millions, millions, et millions d'étoiles !
Je suis, dans l'ombre affreuse et sous les sacrés voiles,
La splendide forêt des constellations.
C'est moi qui suis l'amas des yeux et des rayons,
L'épaisseur inouïe et morne des lumières,
Encor tout débordant des effluves premières,
Mon éclatant abîme est votre source à tous.
Ô les astres d'en bas, je suis si loin de vous
Que mon vaste archipel de splendeurs immobiles,
Que mon tas de soleils n'est, pour vos yeux débiles,
Au fond du ciel, désert lugubre où meurt le bruit,
Qu'un peu de cendre rouge éparsé dans la nuit !
Mais, ô globes rampants et lourds, quelle épouvante
Pour qui pénétrerait dans ma lueur vivante,
Pour qui verrait de près mon nuage vermeil !
Chaque point est un astre et chaque astre un soleil.
Autant d'astres, autant d'immensités étranges,
Diverses, s'approchant des démons ou des anges,
Dont les planètes font autant de nations ;
Un groupe d'univers, en proie aux passions,
Tourne autour de chacun de mes soleils de flammes ;
Dans chaque humanité sont des coeurs et des âmes,
Miroirs profonds ouverts à l'oeil universel,
Dans chaque coeur l'amour, dans chaque âme le ciel !
Tout cela naît, meurt, croît, décroît, se multiplie.
La lumière en regorge et l'ombre en est remplie.
Dans le gouffre sous moi, de mon aube éblouis,
Globes, grains de lumière au loin épanouis,
Toi, zodiaque, vous, comètes éperdues,
Tremblants, vous traversez les blêmes étendues,
Et vos bruits sont pareils à de vagues clairs,
Et j'ai plus de soleils que vous de moucherons.
Mon immensité vit, radieuse et féconde.
J'ignore par moments si le reste du monde,
Errant dans quelque coin du morne firmament,
Ne s'évanouit pas dans mon rayonnement.

Victor Hugo, Extrait du poème "Abîmes" (1853).

Introduction

Contexte historique

La première référence à la possible existence d'objets célestes extérieurs à la Voie Lactée peut être attribuée à l'astronome anglais Thomas Wright dans son ouvrage "On Original Theory or New Hypothesis of the Universe" publié en 1750, et faisant référence à des nébuleuses fruits d'une "création externe". Le concept même de galaxies en tant qu'Univers-îles sera pleinement conceptualisé en 1755 par Emmanuel Kant dans son "Histoire générale et théorie du ciel", très inspirée des principes physiques de mécanique énoncés par Newton. Kant suggère que ces nébuleuses, dont on ne peut distinguer les étoiles qui les composent, se trouvent à des distances infinies, du fait de leur très faible luminosité, et seraient en tous points comparables à la Voie Lactée. Bien que les observations soient peu nombreuses et peu documentées à l'époque, Kant propose une géométrie en disque pour ces nébuleuses particulières, en faisant l'analogie au plan écliptique du système solaire. Au-delà de cette intuition déjà déconcertante d'exactitude, si l'on fait abstraction de la création divine initiale de l'Univers qu'il postule, Kant énonce divers principes cosmologiques qui se sont révélés être très modernes. La cosmologie Kantienne propose un Univers infini qui se trouve dans un état d'évolution perpétuelle. Du chaos initial, émergent des sur-densités de matière qui vont structurer l'Univers grâce aux deux mouvements propres à la mécanique gravitationnelle : la condensation de la matière d'une part et la rotation de la matière d'autre part. Bien que s'inspirant de la mécanique Newtonienne, cette théorie n'est à l'époque ni étayée par les observations, ni par un formalisme mathématique et est considérée, de ce fait, comme un concept plus métaphysique que physique.

Grâce au développement considérable des moyens instrumentaux au début du 20^{ème} siècle, les observations des nébuleuses, ces objets lumineux diffus dont l'origine était encore mal comprise, s'accumulent. Un grand débat fait alors rage dès 1920 dans la communauté des astronomes quant à la nature des nébuleuses spirales. Le débat sur la taille de l'Univers, est principalement animé par deux personnalités qui opposent deux visions radicalement différentes. Harlow Shapley soutient la thèse d'objets faisant partie intégrante de la Voie Lactée, qui rempliraient l'Univers. Herber Curtis défendait lui le concept Kantien d'Univers-île, en dehors de la Voie Lactée. Ce n'est finalement qu'en 1925 qu'Edwin Hubble coupa court au débat en apportant la preuve observationnelle de la nature extragalactique des nébuleuses. En effet, grâce aux observations¹ d'étoiles de type Céphéides dans plusieurs nébuleuses spirales (Andromède, Triangulum, etc.), Hubble met un point final au débat des Univers-îles en prouvant que ces nébuleuses sont non seulement en dehors de la Galaxie, mais qu'il

1. E. Hubble utilisa le télescope Hooker, du nom du mécène aillant financé sa construction, et dont la construction a été achevée en 1917. Il est resté le plus grand télescope du monde jusqu'en 1948 grâce au diamètre de 2.5 mètres de son miroir primaire. Il se situe sur le Mont Wilson en Californie, dans le Comté de Los Angeles aux États-Unis.

s'agissait en outre de galaxies de la même nature que la Voie Lactée. La combinaison des observations de Hubble avec la mesure de décalages spectraux amène un résultat tout aussi important pour la cosmologie moderne, et tout aussi philosophiquement intrigant : les galaxies les plus éloignées présentent un décalage spectral plus important, i.e. une vitesse de récession plus importante. Cette loi nous indique que toutes les galaxies s'éloignent les unes des autres et que les galaxies les plus lointaines sont aussi celles qui s'éloignent le plus vite. Mesurer le décalage spectral des galaxies revient donc à mesurer leur distance par rapport à l'observateur : on parle de "décalage vers le rouge des spectres". Le décalage spectral² z d'une galaxie est exprimé par la formule

$$z = \frac{\lambda_m - \lambda_0}{\lambda_0} = \frac{v}{c} \quad (1)$$

où λ_m est la longueur d'onde décalée par effet Doppler-Fizeau, λ_0 est la longueur d'onde mesurée dans un référentiel au repos, c est la vitesse de lumière et v est la vitesse de récession de l'objet observé. En mesurant le décalage spectral d'un objet suffisamment lointain pour que ses mouvements propres soient négligeables devant la vitesse de récession, on accède donc à l'information de distance de la galaxie.

Bien que le concept d'Univers non statique était quelque peu suspecté à l'époque, la théorie n'était étayée par aucune observation. Il se trouve que le support théorique de cette observation viendra de l'un de ses détracteurs les plus connus : Albert Einstein. En effet, en 1917, Einstein modifie sa théorie de la relativité générale peu après sa publication pour introduire le concept de constante cosmologique. Cette constante, qui peut se traduire en terme d'énergie sous-jacente, permet à Einstein de produire des modèles d'Univers statiques sans commencement ni fin, modèles qui à l'époque, étaient jugés philosophiquement plus acceptables. D'ailleurs, à part le mouvement dans le ciel des planètes (qui signifie étymologiquement astres errant) autres comètes, aucune observation n'indiquait à l'époque que l'Univers n'était pas immuable. Cependant, en modifiant la valeur de la constante cosmologique, il était possible de produire des modèles d'Univers en expansion, comme l'a formalisé Georges Lemaître en 1927, après la découverte du décalage vers le rouge des galaxies en fonction de la distance par Hubble. Du fait de cette émulation scientifique théorique propice, Hubble suggère dans une publication en 1929 que ses observations sont compatibles avec la solution des équations d'Einstein pour un Univers homogène, isotrope et expansion. Les années 1930 marquent donc l'avènement de la cosmologie observationnelle qui naît en même temps que l'astrophysique extra-galactique. La base de la théorie moderne de la cosmologie repose sur la théorie de la relativité générale, qui à ce jour, est le cadre théorique le plus complet permettant d'expliquer un grand nombre d'observations astronomiques.

En 1933, Fritz Zwicky estime la masse dynamique de l'amas de Coma, grâce au théorème de Viriel appliqué sur des mesures de décalages spectraux. En effet, en mesurant la dispersion de vitesses des galaxies qui composent l'amas, il est possible de déduire la masse dynamique du système en faisant l'hypothèse que ce dernier est en équilibre dynamique. Les calculs de Zwicky estimèrent une masse dynamique dix fois plus élevée que la masse lumineuse du système. Bien que les erreurs de mesure aient été grandes, cette observation a été le premier indice de la présence de matière invisible en quantité considérable régissant la dynamique des amas de galaxies. Zwicky est le premier à faire référence à cette masse manquante par la dénomination "matière sombre" dans le sens de son acceptation moderne. Zwicky estima que cette matière sombre pouvait être distribuée dans le milieu

2. "Redshift" en anglais.

intergalactique de l'amas ou bien au contraire être liée aux galaxies individuelles. Jan Oort employa cependant le terme de matière sombre un an avant Zwicky en étudiant le mouvement des étoiles perpendiculairement au plan du disque de la Voie Lactée mais estima que cette masse devait être de la masse baryonique non détectée.

Les galaxies spirales possèdent typiquement une distribution radiale d'étoiles qui décroît avec la distance au centre de la galaxie de manière exponentielle, ce qui produit une courbe de rotation décroissante dans les régions extérieures. La première courbe de rotation a probablement été réalisée par Horace Babcock en 1939 pour la galaxie spirale la plus proche de la nôtre, à savoir M31 (Andromède) et il n'y avait aucune trace de décroissance képlérienne mais Babcock n'a pas considéré l'hypothèse de l'existence de matière noire pour interpréter ces données. Il fallu attendre Mayall en 1951 pour que fut publiée une nouvelle courbe de rotation de la galaxie d'Andromède. Cette dernière ne présentait pas plus de décroissance képlérienne que celle de Babcock 12 ans plus tôt et l'hypothèse de la présence d'une matière noire pour expliquer ces observations ne fut sérieusement avancée. D'autres galaxies proches comme M33 ont été réalisées avant la fin des années 1960, par exemple par des astronomes de l'Observatoire de Marseille (e.g. Dumoulin, Viton & Georgelin, 1968) mais le dogme de la décroissance attendue n'a pas permis à tous ces observateurs d'interpréter les données avec une vision différente. L'émergence de la radio astronomie permit d'explorer la dynamique des galaxies en utilisant le gaz neutre comme traceur (la raie à 21 cm) plutôt que le gaz ionisé comme c'est le cas dans le domaine optique, limitant les observations aux régions internes des galaxies. La première courbe de rotation étendue au-delà du disque stellaire optique de M31 fut publiée en 1957 par van der Hulst et collaborateurs. Ils montrèrent que la vitesse de rotation ne décroissaient pas sensiblement avec la distance au centre de la galaxie, très éloignée du comportement attendu pour une décroissance képlérienne sans que cela ne déclenche véritablement un sentiment remise en cause dans la communauté. Il fallu véritablement attendre la fin des années 1970 pour que, dans le cadre de la dynamique newtonienne, la nécessité de faire appel à de la matière noire pour interpréter les courbes de rotation soit établie. L'obtention de données de meilleure qualité ayant permis une interprétation plus précise et les progrès théoriques sur la compréhension de la dynamique des disques, ont conduit l'ensemble de la communauté à accepter que l'existence de courbes de rotation globalement plates dans la totalité des galaxies spirales est en complète contradiction avec la dynamique newtonienne, si l'on suppose que seules les étoiles et le gaz contribuent au potentiel gravitationnel de la galaxie. Parallèlement, dans les années 1980, les succès de la cosmologie et l'émergence d'une nouvelle discipline, celle des astroparticules, ont fait émerger l'idée que cette matière noire n'était pas baryonique. Des modifications à la théorie classique offrent cependant un cadre intéressant, mais pas encore une théorie complète, pour interpréter les courbes de rotation plates, ainsi que d'autres manifestations supposées de la matière noire, sans faire appel à cette dernière, la plus célèbre des alternatives étant MOND (Modified Newtonian Dynamics).

En 1964, Arno Penzias et Robert Wilson, deux ingénieurs de l'entreprise de téléphonie américaine Bell, découvrirent un rayonnement radio isotrope qui perturbait leurs mesures. Malgré un effort particulier sur la réduction de toutes les sources de bruits, un signal 100 fois plus important que prévu persistait dans les données. La proximité géographique des ingénieurs avec l'Université de Princeton leur permit d'avoir accès aux travaux encore non publiés de Jim Peebles, Robert Dicke, et David Wilkinson qui prédisaient une émission radio relique dans le cas d'un scénario d'une cosmologie initiée par une explosion primordiale. C'est à la lumière de ces travaux que les deux ingénieurs se rendirent compte de l'importance considérable de leur mesures. La prédiction théorique de ce rayonnement cosmologique

comme conséquence directe de la théorie du Big Bang avait d'ailleurs été émise en 1948 par Alpher, Bethe & Gamow. La découverte observationnelle fortuite de ce rayonnement permit à Penzias et Wilson d'être récompensés du prix Nobel de physique en 1978, et pour cause : l'observation de ce signal donne une base observationnelle solide à la théorie du "Big Bang". Ce signal est interprété comme étant une signature directe de la radiation thermique émise par l'Univers au moment de sa ré-ionisation.

Cosmologie moderne

Les trois piliers classiques de la cosmologie moderne sont :

1. L'abondance des éléments primordiaux, qui permet d'expliquer la présence des noyaux légers (H, D, T, He, Li, Be, B, ...) lors de la nucléosynthèse primordiale, également initiée par Gamow,
2. L'expansion de l'Univers (loi de Hubble)
3. L'existence du fond diffus cosmologique.

Après que les neutrinos aient occupés le devant de la scène en tant que candidat de matière noire et chaude non baryonique, cette hypothèse fut finalement abandonnée. Le scénario cosmologique dans lequel la matière noire froide (Blumenthal et al., 1984) domine la distribution de masse de l'Univers est aujourd'hui devenu un modèle "standard" accepté par dans la communauté scientifique. Dans ce modèle, les particules qui composent la matière sombre sont très massives et se déplacent lentement. C'est cette propriété qui favoriserait la formation de galaxies en tant que structures primordiales de l'Univers, et qui par la suite, se regrouperaient dans des amas eux-mêmes se regroupant dans des super-amas : on parle alors de scénario de formation des structures "du bas vers le haut"³. Ce scénario est aujourd'hui largement accepté notamment grâce à l'aptitude des simulations numériques incluant un modèle de matière noire froide à reproduire de nombreuses propriétés physiques des galaxies et des structures à grandes échelles de l'Univers (voir figure 0.2).

Le modèle de concordance de la cosmologie moderne qui a émergé dans les années 1990 est le modèle dit " Λ CDM" pour "Lambda Cold Dark Matter". Le terme "lambda" fait référence à la constante cosmologique, dont la valeur positive traduit l'existence d'une énergie du vide. À ce jour, il est considéré comme étant le meilleur modèle de Big Bang dont nous disposons. Il existe une grande variété de sous-modèles avec des raffinements plus ou moins exotiques, mais de manière générale, un modèle qui tient compte de la matière noire et de l'énergie noire est considéré comme un modèle Λ CDM. Ce modèle permet d'expliquer avec une bonne précision les trois piliers fondamentaux de la cosmologie moderne ainsi que bon nombres d'observations fondamentales plus récente de la cosmologie observationnelle, à savoir :

1. La découverte des fluctuations de température dans le fond diffus cosmologique grâce au satellite COBE en 1992 qui ont probablement donné naissance aux structures à grande échelle observées aujourd'hui dans l'Univers. C'est la naissance de la cosmologie dite de précision.
2. L'observation des structures à grandes échelles de la distribution des galaxies et des amas de galaxies dans l'Univers proches. Ces observations ont été réalisées à partir des grands sondage spectroscopiques de galaxies

3. "downsizing" en anglais.

3. L'observation des pics acoustiques et donc des petites échelles de l'Univers dans l'analyse statistique des anisotropies du fond diffus cosmologique. Ces oscillations acoustiques baryoniques prédites dès 1970 ont été observées pour la première fois en 1999 (expérience ballon Boomerang), largement confirmé ultérieurement par le satellite WMAP (2007).
4. La nécessité de faire appel à de la matière noire dans les amas de galaxies pour expliquer leur masse. Ces mesures sont réalisées grâce aux observations du gaz très chaud qui émet en rayon X au centre des amas, à l'aide de l'effet de lentilles gravitationnelles que l'amas induit sur l'image d'une source d'arrière plan à l'amas, ou encore des observations spectrales optiques permettant de mesurer la dispersion de vitesses des galaxies. La découverte d'une collision entre deux amas que constitue l'amas du boulet a apporté récemment (2006) une contrainte forte sur la distribution et l'existence de matière non baryonique au sein des deux amas.
5. L'accélération de l'expansion de l'Univers telle que dérivée de la relation redshift-distance des sondes cosmologique (supernovae de type Ia, 1998, 1999) et d'autres indicateurs nécessitant de faire appel à de l'énergie noire.

Ces nombreuses mesures observationnelles permettent aujourd'hui de calculer les paramètres cosmologiques décrivant l'Univers, dans le cadre de la théorie de la relativité générale et d'obtenir de fortes contraintes sur le modèle standard. En se référant à la première analyse des données collectées du fond diffus cosmologique par l'équipe du télescope spatial Planck en mars 2013 (voir figure 0.1), la densité énergétique de l'Univers est partagée entre l'énergie noire ($\Omega_\Lambda=0.692\pm 0.010$), la matière noire ($\Omega_c=0.1187\pm 0.0017$), et la matière baryonique ($\Omega_b=0.02214\pm 0.00024$), avec une constante de Hubble estimée à $H_0=67.80\pm 0.77$ km.s⁻¹.Mpc⁻¹. Ces paramètres permettent d'estimer l'âge de l'Univers à 13.798 ± 0.037 milliards d'années.

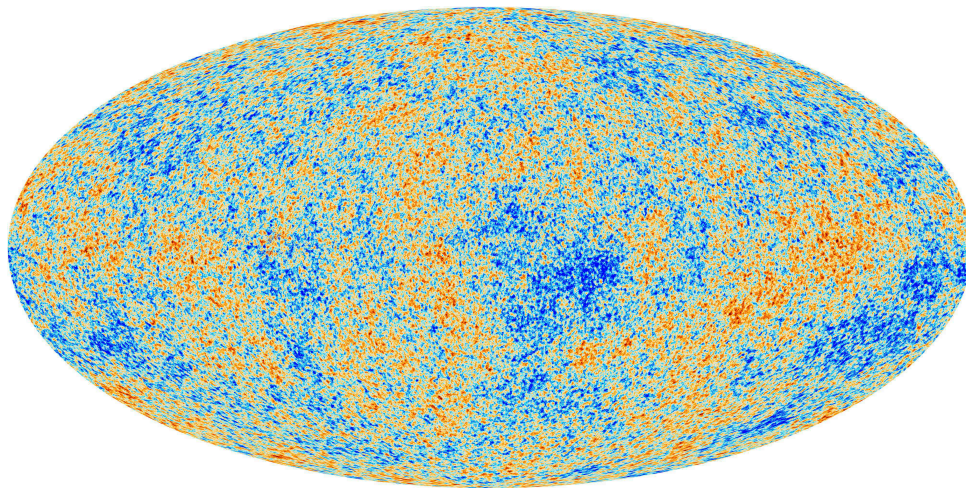


Figure 0.1 – Le fond diffus cosmologique vu par le satellite Planck. Crédits : ESA.

Dans ce modèle, la formation des structures cosmologiques est provoquée par des perturbations quantiques infimes du champ de densité primordial de la matière qui sont propagées à des échelles cosmologiques par la période d'inflation suivant le Big Bang. L'effondrement gravitationnel de la matière noire vers ces zones de sur-densités permet une croissance initialement linéaire de ces perturbations, qui est modélisée par la théorie des systèmes linéaires. C'est lorsque ces perturbations deviennent non-linéaires que la description de la

croissance de ces perturbations devient sensiblement plus compliquée. Les simulations N-corps cosmologiques deviennent alors un outil numérique de choix permettant d'explorer le comportement de la matière noire dans le régime non-linéaire. De nombreux travaux tant sur le plan analytique que sur le plan numérique montrent la formation d'une superstructure composée de halos de matière sombre reliés entre eux par des filaments. Cette organisation de la matière sur les plus grandes échelles de l'Univers est plus communément désigné par le terme de "toile cosmologique". Cette structure en filaments est confirmée par les grands relevés cosmologiques (voir figure 0.2), et semble être la plus grande échelle d'organisation de la matière dans l'Univers.

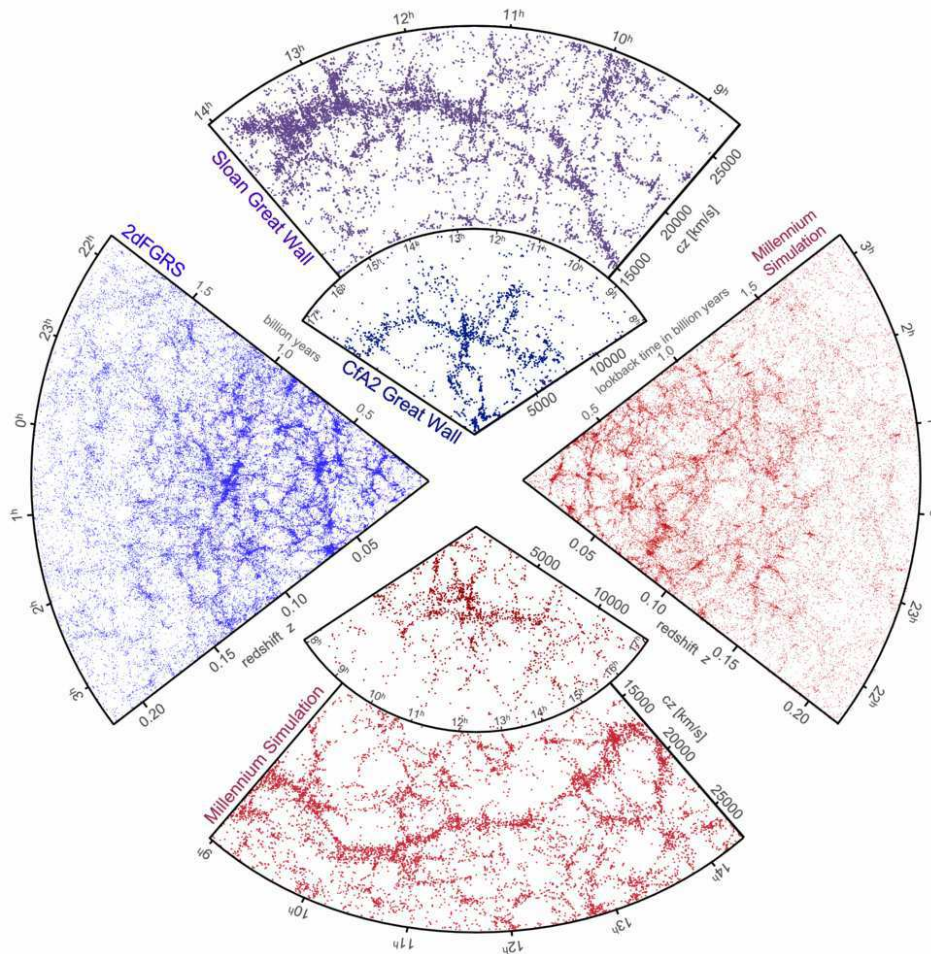


Figure 0.2 – Comparaison des structures à grandes échelles formées par les galaxies provenant de grands relevés spectroscopiques ainsi que de simulations cosmologiques. Les quadrants de gauche et du haut (couleur bleue) exposent la distribution des distances radiales des galaxies de notre voisinage observées dans différents grands relevés (2dFGRS, SDSS, CfA2). Les quadrants de droite et du bas (couleur rouge) montrent la distribution des distances radiales de galaxies issues d'un catalogue comparable aux relevés observationnels et construit grâce à la simulation Millenium. Crédits : Springel et al. (2006).

Le résultat de la croissance non-linéaire d'une perturbation du champ de densité de la matière sombre est une structure désignée par le terme de "halo". Il s'agit d'un objet stable, en équilibre dynamique grâce aux mouvements aléatoires des particules qui le composent. Dans le modèle Λ CDM, la matière baryonique et non-baryonique sont initialement distri-

buées de manière isotrope et homogène, suivant la distribution de matière noire sur des échelles supérieures à la longueur de Jeans locale. La conservation du moment angulaire lors de l'effondrement gravitationnel des baryons mène à la formation d'une structure en disque dont la rotation contrebalance la gravitation. Le même phénomène se produit pour la matière noire, mais à la différence de celle-ci, le gaz est une composante dissipative qui est capable de rayonner son énergie interne, et donc de se refroidir. De manière générale, on estime que le moment angulaire spécifique du disque est égal au moment angulaire spécifique du halo de matière sombre, bien qu'il puisse exister un désalignement entre les vecteurs de moment angulaire du halo et du disque baryonique dans les simulations numériques. La première génération de halo est le résultat de l'évolution de ces perturbations primordiales, mais l'interaction gravitationnelle plus tardive entre les halos peut donner lieu à des transferts de moment angulaire orbital, et provoquer plus particulièrement des fusions qui sont censées jouer un rôle majeur dans l'évolution de leur propriétés physique au cours des différentes ères cosmologiques.

La finitude de la vitesse de propagation des photons est une chance pour les astronomes, elle permet l'observation des objets les lointains et donc de remonter dans le temps et d'étudier la formation et l'évolution des briques de l'Univers que sont les galaxies au travers des âges.

Structure des galaxies

Une galaxie est un amas d'étoiles, de gaz, de poussières, et de matière noire. Les galaxies de l'Univers local offrent une grande variété de types morphologiques, de tailles et de masses. Les plus ténues d'entre elles contiennent quelques 10^7 étoiles, tandis que les galaxies les plus massives abritent plus de 10^{12} étoiles. Non content d'avoir établi la nature extra-galactique de ces nébuleuses, Edwin Hubble a également mis au point (1924-1936) une classification des types morphologiques des galaxies locales (voir figure 0.3). On se réfère généralement à cette classification par le terme "séquence de Hubble". On trouve dans cette séquence :

- Les galaxies spirales : ce sont les galaxies avec une géométrie planaire, que l'on peut décomposer en un disque en rotation autour d'un bulbe stellaire. Une fraction importante de ces galaxies présentent une structure spirale dans le disque, dont le degré d'enroulement varie d'une galaxie à l'autre. Une explication de la fréquence d'observation de structures spirales dans les disques serait l'existence d'ondes de compression se propageant dans les disques (Lin et Shu, 1964). Dans les régions qui constituent la structure spirale, on trouve un gaz comprimé plus dense dont la formation stellaire est supérieure au reste du disque. La plupart de ces galaxies spirales sont dites "barrées", du fait de la structure rectiligne qui traverse le noyau galactique. Le disque est le siège de la formation stellaire de la galaxie, où les étoiles jeunes rayonnent leur énergie dans les longueurs d'ondes courtes du domaine visible (bleu) tandis que le bulbe est lui composé d'étoiles vieilles qui auraient migré depuis le disque vers des orbites internes plus excentriques.
- Les galaxies elliptiques : ce sont les galaxies dont la forme évoque effectivement un ellipsoïde plus ou moins aplati. Les propriétés physiques de ces galaxies les rendent bien différentes des galaxies spirales. Contrairement aux galaxies spirales, les galaxies elliptiques forment peu ou pas d'étoiles du fait de très faibles quantités de gaz. Leurs

étoiles sont vieilles, et le support dynamique est assuré par les mouvements aléatoires de leurs étoiles. On les oppose aux galaxies spirales bleues du fait que leurs vieilles étoiles soient d'une couleur rouge caractéristique. Les galaxies elliptiques, comme les galaxies spirales d'ailleurs, peuvent abriter un trou noir supermassif au centre de masse.

- Les galaxies lenticulaires : ce sont des galaxies intermédiaires, à mi-chemin entre les galaxies spirales et les galaxies elliptiques. Elles se caractérisent par un grand bulbe massif, autour duquel gravite un disque peu actif en terme de formation stellaire et où l'on ne distingue pas de structures spirales.
- Les galaxies irrégulières : ce sont les galaxies qui ne rentrent pas dans les classes définies ci-dessus. Elles exhibent des morphologies chaotiques, composées d'étoiles jeunes et de grandes quantités de gaz.

Au delà de la simple classification morphologique, Hubble voyait là une véritable séquence évolutive, où les galaxies de type précoce (galaxies elliptiques) évoluent au cours du temps pour former les galaxies du type tardif (galaxies spirales). Bien que l'intuition de Hubble concernant la capacité des galaxies à modifier leur morphologies au travers des âges était juste, le sens de l'évolution était lui partiellement faux. On sait aujourd'hui que l'évolution se produit plutôt dans le sens inverse, i.e. des galaxies spirales vers les galaxies elliptiques par l'intermédiaire des fusions, du moins pour former les galaxies elliptiques les plus massives.

Les galaxies de l'Univers distant montrent des morphologies plus complexes et la question des processus physiques conduisant à la formation de la séquence de Hubble demeure une interrogation fondamentale de l'astrophysique moderne. L'observation de champs profonds, comme par exemple les champs profonds et ultra profond du télescope Hubble réalisées en 1995, 1998 et en 2004 ont suggéré que l'Univers contenait des centaines de milliards de galaxies mais aussi que ces galaxies lointaines et conséquemment jeunes étaient morphologiquement très différentes des galaxies actuelles. Tout d'abord la proportion de galaxies irrégulières ainsi que de galaxies perturbées était très élevée par rapport aux proportions actuelles et la fraction de galaxies elliptiques géantes faible. D'autre part, les galaxies étaient aussi globalement plus petites, plus gazeuses et plus chaotiques que maintenant. Ces observations signifient probablement que les grands disques stables ainsi que les galaxies elliptiques géantes que nous observons aujourd'hui n'étaient pas encore constitués et qu'ils ne se sont formés ultérieurement par la fusion successive de galaxies irrégulières plus petites. L'Univers étant plus dense, les fusions entre galaxies étaient d'ailleurs plus fréquentes que dans l'Univers actuel. Enfin, ces galaxies étaient aussi plus bleues, ce qui signifie qu'elles formaient plus d'étoiles que maintenant. La variété des galaxies observées à différents décalages spectraux a permis d'estimer la variation du taux de formation stellaire en fonction de l'âge de l'Univers et de montrer que cette formation a culminé il y a environ 10 milliards d'années. Le système de classification de Hubble qui a pour objectif de classer les galaxies proches ne peut pas s'appliquer aux galaxies lointaines. Par contre, l'évolution des galaxies doit conduire à la classification de Hubble et pas seulement à la formation des galaxies elliptiques géantes, il est donc nécessaire que la formation des galaxies primordiales et que leur évolution ultérieure par fusions successives de galaxies, par accrétion continue de gaz froid intergalactique et par évolution séculaire conduisent non seulement à la constitution des galaxies elliptiques géantes parfaitement relaxées mais aussi à celles des grands disques réguliers observés aujourd'hui. Delgado-Serrano et al. (2010) ont tenté une classi-

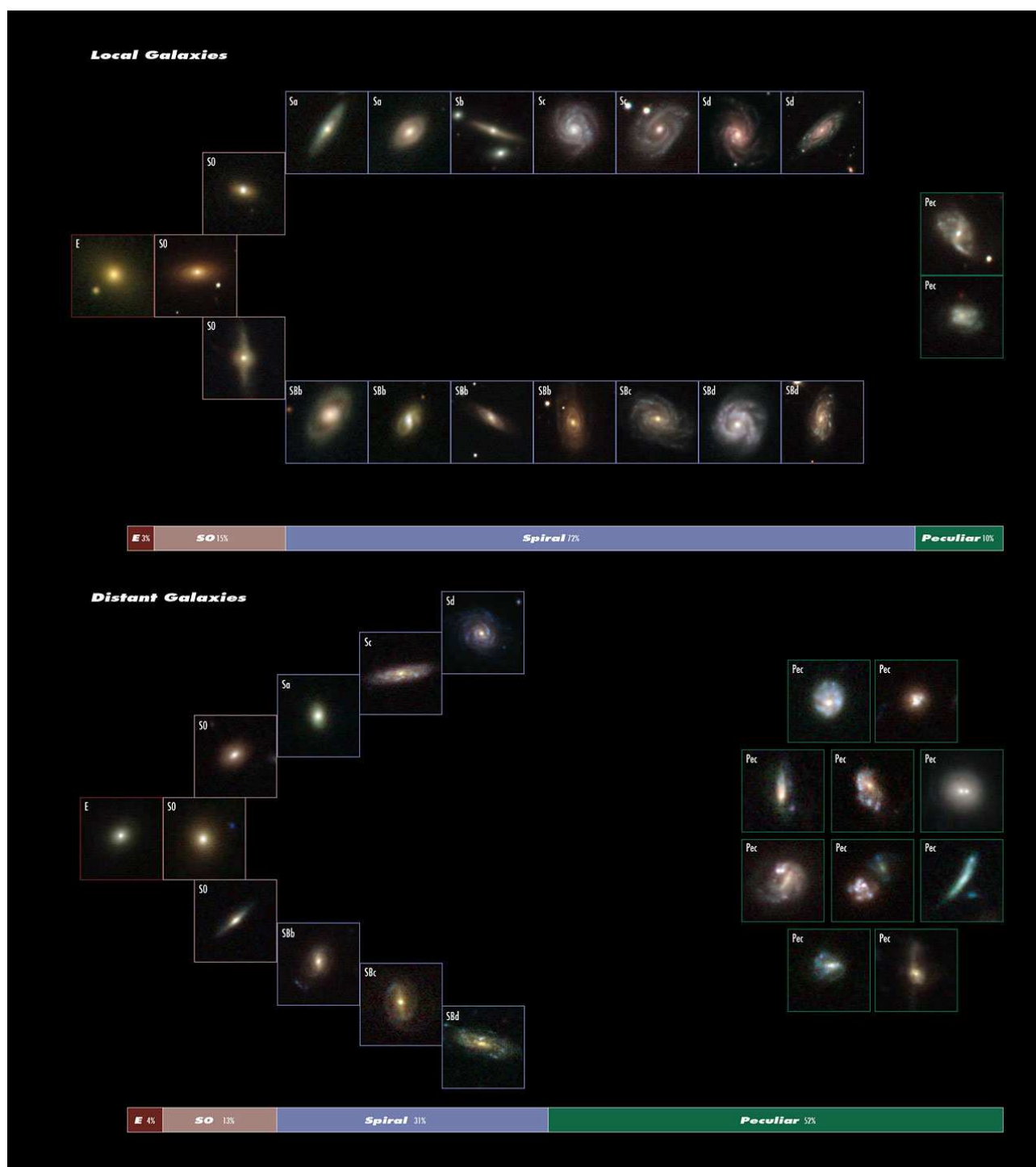


Figure 0.3 – Schémas de classification des galaxies, différenciant les galaxies elliptiques (à gauche), spirales (au centre) et irrégulières ou “péculières” (à droite). Les galaxies elliptiques sont désignées par la lettre “E”, tandis que les galaxies spirales sont désignées par la lettre “S”. Les galaxies spirales barrées sont désignée par le suffixe “B”. Ces figures ont été créées à partir de données prises par le télescope du “Sloan Digital Sky Survey” (figure du haut) et du télescope Hubble (figure du bas), elles permettent de comparer les galaxies actuelles aux galaxies environ 6 milliards d’années plus jeunes. La figure du haut illustre la classification de Hubble en forme de diapason qui représente les galaxies actuelles. La figure du bas montrent une fraction beaucoup plus importante de galaxies irrégulières. Les fractions respectives de type morphologique aux deux différentes époques sont représentées par la barre colorée sous les deux images. Crédits : Delgado-Serrano et al. (2010)

fication des galaxies à grand décalage vers le rouge qui illustre bien ce propos (voir figure 0.3).

Formation stellaire

Les galaxies contiennent une partie de leur masse sous forme de gaz, essentiellement de l'hydrogène (~76%) et de l'hélium (~23%). Le gaz situé dans le disque est plus communément appelé milieu interstellaire, et constitue un réservoir qui alimente la formation stellaire de la galaxie tout au long de son existence. La formation des étoiles à partir d'une masse de gaz en effondrement gravitationnel est un phénomène physique extrêmement complexe qui est une théorie à part entière. À l'échelle galactique, on a depuis longtemps résumé la formation stellaire à une loi de Schmidt (1959) et Kennicutt (1998), qui stipule que le taux de formation stellaire par unité de surface $\dot{\Sigma}_\star$ est directement proportionnel à la densité surfacique de gaz Σ_{gaz} :

$$\dot{\Sigma}_\star \propto \Sigma_{gaz}^n \quad (2)$$

avec n une constante empirique estimée à 1.4 dans les galaxies qui forment encore des étoiles de façon conséquente dans l'Univers local. Cette loi purement empirique souffre évidemment de tous les défauts associés aux déductions sans réelle consistance physique, mais permet d'appréhender les quantités d'étoiles formées par une grande variété de galaxies. Les observations de régions de formation stellaire dans la Voie Lactée montrent que les étoiles se forment dans le gaz sous forme moléculaire, où les densités dépassent 1000 atomes d'hydrogène par cm^{-3} . La densité de ce gaz est telle qu'il devient opaque au rayonnement des étoiles environnantes, et permet donc d'avoir un coeur extrêmement froid et dense où vont se former les étoiles. Une grande fraction du gaz dans l'environnement proche des étoiles de type spectral OB (étoiles jeunes et dont les masses sont supérieures à dix masses solaires) absorbe les rayonnements fortement ionisants de ces étoiles. On désigne le gaz ionisé autour de ces étoiles par l'appellation "région HII", du nom de la transition électronique atomique permettant au gaz de ré-émettre des photons essentiellement à 13.6 électrons-volts. Cette émission n'est malheureusement pas un bon traceur des régions de formation stellaires car ce rayonnement est absorbé aussitôt émis dans les milieux optiquement épais mais il existe d'autres raies d'émission, comme la raie $H\alpha$, qui est généralement suffisamment intense et qui permet de mesurer cette formation stellaire dans toutes les galaxies de l'Univers qui forment des étoiles. Lorsqu'elles atteignent la fin de leur vie, au bout de quelques millions d'années, les étoiles de type OB s'effondrent sur elles-mêmes et atteignent des densités dans leur coeur telles que les réactions de fusion thermonucléaire ne sont plus en mesure de contrebalancer le poids de l'astre. Des réactions de fusions se propagent jusqu'aux couches supérieures qui sont expulsées en quelques secondes seulement. Les explosions de ces astres en fin de vie (ou supernovae) associée à ces processus relâchent des quantités phénoménales d'énergie dans leur milieu environnant. La puissance dégagée par les supernovae est si importante qu'on peut les observer dans des galaxies distantes de plusieurs milliards d'années lumière. Bien qu'il existe différents processus menant à cette réaction, nous nous intéressons ici à l'effet des supernovae à l'échelle galactique. Les étoiles les plus massives ont des durées de vie courtes comparées aux échelles de temps cosmologiques, et injectent donc en permanence de l'énergie et du gaz dans le milieu interstellaire. Dans les simulations de galaxies, on modélise la contribution des supernovae au bilan énergétique total de la galaxie en injectant une énergie sous forme thermique et/ou cinétique afin de modéliser la propagation de l'onde

de Sedov générée par la réaction thermonucléaire. Les supernovae relâchent une partie des baryons qui la constituait sous forme de gaz. Ce gaz est enrichi par les métaux, produits de la fusion nucléaire, et se dissout dans le milieu interstellaire environnant, participant ainsi à la formation de la prochaine génération d'étoiles.

Support dynamique des galaxies

Comme nous l'avons vu au cours de la description de la séquence de Hubble, le support dynamique des galaxies n'est pas le même que l'on parle d'une galaxie spirale où bien d'une galaxie elliptique. Les étoiles sont une composante non-dissipative (car il s'agit d'un système non-collisionnel) des galaxies. Cela signifie que ces dernières n'ont aucun moyen de se refroidir une fois chauffées. Bien sûr, le seul processus de chauffage dont nous pouvons parler dans le cas du système N-corps est le chauffage gravitationnel induit par la friction dynamique et les forces des marées, qui peuvent provenir aussi bien de processus d'évolution séculaires (formation d'une barre), ou bien d'interactions avec l'environnement (interactions et/ou fusions avec des galaxies environnantes). Le gaz, quant à lui, est une composante dissipative : son rayonnement, en particulier celui des raies d'émission des métaux permet au gaz se refroidir. Ce refroidissement s'accompagne de la création d'un disque qui est le résultat naturel de la conservation du moment angulaire.

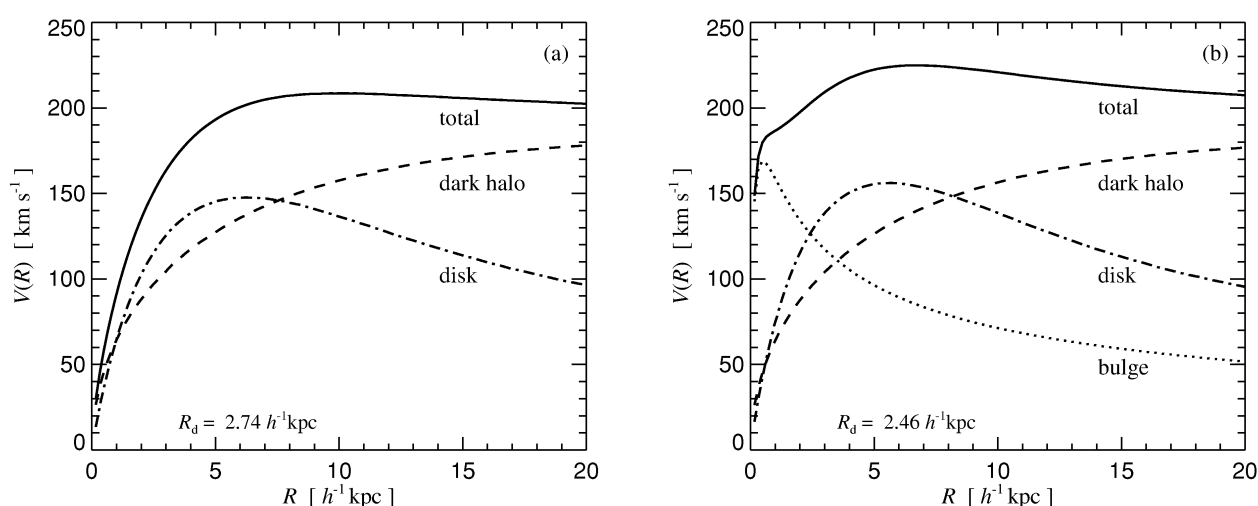


Figure 0.4 – Courbe de rotation pour un modèle de galaxie avec bulbe stellaire (à droite), et sans bulbe stellaire (à gauche). Crédits : Springel et al. (2005b)

Comme nous l'avons évoqué dans la section décrivant le contexte historique de l'introduction de la matière noire pour expliquer les courbes de rotation des galaxies, on caractérise le support dynamique d'un disque en mesurant sa courbe de rotation, qui trace la vitesse circulaire en fonction de la distance au centre galactique. Dans le cas d'un disque à l'équilibre dynamique, (i.e. dont la distribution spatiale de masse ne varie pas sur plusieurs centaines de millions d'années, soit plusieurs rotations galactiques), la vitesse circulaire est une mesure directe du potentiel gravitationnel de la galaxie. Chacun des éléments qui la constituent (étoiles dans le disque et dans le bulbe, gaz dans le disque et matière noire dans le halo) contribue à une fraction de la courbe de rotation (voir figure 0.4). La capacité de déterminer observationnellement la courbe de rotation d'une galaxie par effet Doppler sur les raies d'émission ou d'absorption est d'ailleurs l'une des explications à l'origine du concept

de matière noire : la distribution exponentielle de la lumière dans les galaxies suggère une courbe de rotation avec un pic situé à un rayon d'environ deux fois la longueur d'échelle du disque, puis une décroissance prononcée dans les parties externes du disque. Or, la plupart des courbes de rotation des galaxies de l'Univers local sont plates. Le moyen de concilier la théorie de la gravitation et les propriétés dynamiques des galaxies était donc de supposer l'existence d'un halo de matière non-visible contribuant au potentiel gravitationnel et compensant la décroissance de la courbe de rotation dans les régions externes du disque.

Interaction entre galaxies

L'interaction et la fusion de galaxies locales est un processus connu depuis déjà de nombreuses années du fait que l'on observe une fraction significative de galaxies en paire avec des signatures morphologiques et cinématiques qui traduisent la présence de forces de marées (voir figure 0.5). La fusion est un processus évolutif dont il est impossible de comprendre la nature seulement à partir des observations. Ce sont les simulations numériques qui ont permis de véritablement cerner les processus menant à coalescence complète de deux galaxies. Le travail pionnier de Toomre et Toomre (1972) a mis en évidence le lien entre queues de marée et interactions galactiques.



Figure 0.5 – Paire de galaxies spirales de masse équivalente en interaction (Arp 240). Crédits : NASA, ESA, The Hubble Heritage - ESA/Hubble collaboration, et A. Evans (University of Virginia, Charlottesville/NRAO/Stony Brook University).

Le premier passage au périastre de la trajectoire Keplerienne de chacune des deux galaxies qui se rencontrent déclenche d'intenses forces de marée, qui projettent de grandes quantités de gaz et d'étoiles dans de longues structures allongées désignées par le terme

“queues de marée”. Le processus de friction dynamique est à l’origine du déclin orbital, transférant une partie du moment angulaire orbital vers l’énergie interne du halo de matière noire. Une partie des grandes quantités de gaz éjecté peut ensuite retomber sur la galaxie et nourrir alors une flambée de formation stellaire très concentrée dans le noyau galactique. Ce schéma est cependant bien plus compliqué qu’il ne le laisse entendre : de nombreux paramètres très différents d’un cas à l’autre contrôlent la vitesse de fusion, la formation stellaire induite, la morphologie du rémanent, etc. De manière générale, les simulations numériques montrent que la fusion des deux galaxies spirales ont tendance à former une galaxie elliptique, prenant à rebours la séquence évolutive décrite par Hubble. La présence de grands réservoirs de gaz est néanmoins capable de reconstruire les disques grâce au refroidissement rapide de ce gaz dans le plan du disque. La fusion de galaxie est donc un processus important dans l’évolution d’une galaxie, qui accélère sa transition morphologique, dynamique et chimique. Dans les modèles cosmologiques, on évalue la contribution des fusions à l’assemblage de masse de galaxies à environ un tiers de la masse totale accrétée par une galaxie au cours de sa vie (Dekel et al., 2009a). La validation de tels modèles cosmologiques passe donc par la contrainte observationnelle de l’occurrence statistique des fusions à tous les moments de l’histoire de l’Univers.

Plan de la thèse

La problématique dans laquelle s'inscrit cette thèse peut se résumer aux questions suivantes : qu'est-ce qui caractérise les fusions de galaxies jeunes très riches en gaz ? Peut-on détecter des signatures de fusions récentes par l'analyse de la cinématique d'une galaxie lointaine ? Les différentes parties de ce manuscrit sont présentées de manière à refléter mon travail au cours de cette thèse dans l'ordre chronologique dans lequel il a été réalisé. Cette articulation permet d'appréhender la démarche scientifique progressive dans laquelle ce projet s'est développée. Les différents chapitres s'organisent de la manière suivante :

- Dans un premier chapitre, je présente une nouvelle technique d'analyse cinématique de l'échantillon MASSIV. Pour ce faire, j'ai travaillé au développement d'une méthode automatisée ne faisant plus appel à une classification visuelle. Cette méthode repose essentiellement sur un algorithme génétique couplé à une décomposition en harmoniques sphériques du champ de vitesses radiales (aussi appelé "kinemetry"). Bien que cette méthode soit expérimentée sur des données de galaxies lointaines à faible résolution spatiale, elle peut tout à fait être étendue à des données de galaxies beaucoup plus proches avec une résolution spatiale plus importante. Ce chapitre sera l'occasion de montrer qu'une simple analyse des observations n'est pas suffisante, et qu'il est nécessaire de disposer de modèles numériques détaillés de fusions afin de pouvoir inférer d'éventuelles signatures. Ce chapitre, qui résulte d'une préparation de publication, est présenté en anglais.
- Dans un second chapitre, j'introduis une méthode de construction de conditions initiales de galaxies pour simulations numériques idéalisées. Je présente de manière didactique comment j'ai été en mesure de bâtir des distributions de particules avec des vitesses permettant d'obtenir un modèle numérique de galaxie à l'équilibre dynamique. Cette méthode, que j'ai développée au cours de ma deuxième année de thèse, était un préalable requis pour pouvoir définir un échantillon de simulations numériques permettant de comprendre la dynamique complexe du gaz dans les galaxies à grand décalage spectral. J'ai porté une attention particulière sur la mise à disposition à la communauté scientifique de l'implémentation numérique de la méthode via un logiciel libre d'accès.
- Dans un troisième chapitre, j'introduis l'échantillon de simulations numériques MIRAGE, dont j'ai spécifié les propriétés afin qu'il permette une meilleure compréhension de la cinématique du gaz dans les fusions de galaxies à grand décalage spectral. La définition, la réalisation, et l'analyse de cet échantillon composent la majeure partie de mon travail au cours de ma troisième année de thèse. Cet échantillon s'oriente sur une exploration statistique de deux paramètres majeurs contrôlant les propriétés physique des rémanents de fusion galactique : la masse et l'orientation initiale des deux disques. Cet échantillon fait écho à la problématique de classification cinématique des galaxies des échantillons spectroscopiques lointain. Il a été défini de manière à répondre à la question : peut-on détecter des signatures de fusions dans la cinématique du gaz des galaxies à $z > 1$ dans des conditions observationnelles ? Ce chapitre présente une première publication dans laquelle je définis l'échantillon et je réalise une première analyse détaillant les propriétés globales des simulations.
- Dans un quatrième chapitre, j'aborde la problématique de la présence de granules dans la morphologie des galaxies "juvéniles". Je présente notamment l'analyse des

structures en granules dans les simulations MIRAGE qui sont formées par les instabilités gravitationnelles naturellement présentes dans les disques des galaxies jeunes. En effet, l'échantillon MIRAGE étant focalisé sur l'étude de galaxies à très forte fraction de gaz, l'analyse détaillée des sous-structures formées par effondrement gravitationnel est indispensable car ces structures sont à l'origine de la majeure partie de la formation stellaire dans les galaxies de l'Univers lointain. Notamment, je montre que ces granules sont compatibles avec un scénario où ceux-ci persistent pendant plusieurs centaines de millions d'années jusqu'à compléter leur migration vers le centre de la galaxie hôte. La dernière partie de ce chapitre, qui met en avant les processus d'évolution conduits par les instabilités gravitationnelles, fait l'objet d'une préparation de publication.

- Dans un cinquième et dernier chapitre, je présente les résultats de l'étude préliminaire comparative entre les simulations numériques réalisées dans le cadre de cette thèse et l'échantillon de galaxies lointaines MASSIV. La comparaison est réalisée grâce à la création de plus de 4000 pseudo-observations qui projettent les simulations MIRAGE au décalage spectral $z=1.7$ et simulent la réponse instrumentale du spectrographe SINFONI installé au foyer du VLT. Ce dernier chapitre a vocation à apporter des éléments nouveaux quant à la détection de signatures de fusions de galaxies très riches en gaz dans les échantillons spectroscopiques 3D.

Analyse cinématique de l'échantillon



Résumé

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Dans ce chapitre, rédigé en anglais du fait qu'il a été en partie construit à partir d'un article en préparation, je présente les données de l'échantillon spectroscopique MASSIV composé de 83 galaxies observées avec le spectrographe intégral de champ SINFONI au télescope VLT. Après avoir passé en revue les principaux résultats scientifiques issus de l'analyse de ces observations, j'explique ma contribution au développement d'une méthode d'analyse cinématique basée sur une décomposition en harmoniques sphériques. L'analyse ainsi réalisée est comparée aux précédentes analyses de la littérature.

1.1 MASSIV observational context

The questions related to the evolution processes leading galaxies towards the Hubble sequence can be addressed with spatially resolved measurements from 3D spectral analysis, which allows to probe various physical quantities; e.g. star formation, gas kinematics, chemical abundances. The MASSIV sample investigates a redshift range that was unexplored by Integral Field Spectroscopy (IFS), considering that four distinct 3D surveys are preceding it. The IMAGES sample (Puech et al., 2008), with $0.4 < z < 0.75$, showed that regular rotating disks are quite similar to the local ones and that mergers are playing an important role in galaxy mass assembly in this redshift range. The SINS sample (Förster Schreiber et al., 2009), Wright et al. (2009) and Law et al. (2009) samples, with galaxies in the range $1.5 < z < 3$, and the LSD/AMAZE sample (Maiolino et al., 2008) at larger redshift ($3 < z < 4$), showed that a lot of young galaxies are experiencing a high gaseous turbulence.

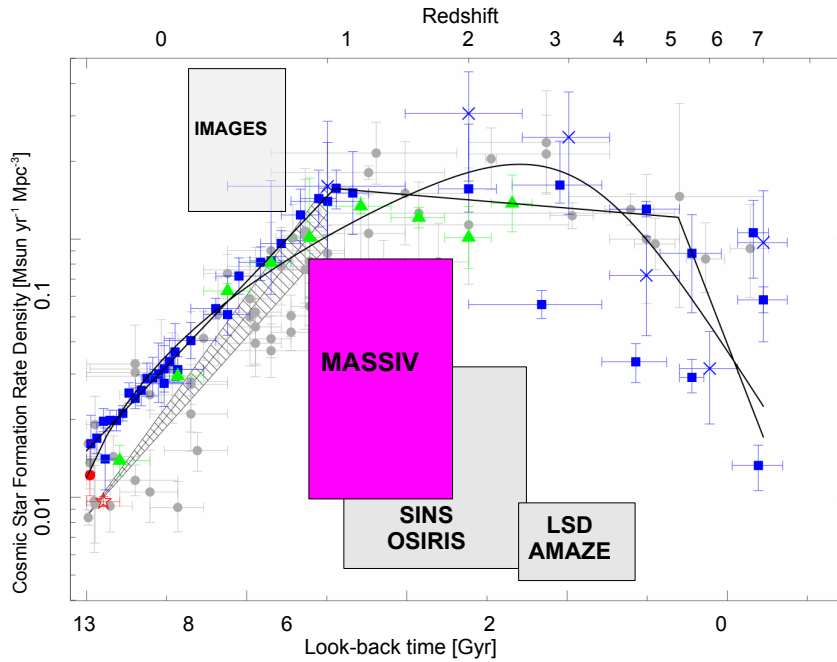


Figure 1.1 – Evolution of the cosmic star formation rate density as a function of look-back time and redshift. Major IFU surveys redshift ranges are compared together. The relative height of each box is proportional to the sample size. Credits : Contini et al. (2012b).

The MASSIV sample was built to probe the galaxies representative of the star-forming population in the redshift range $0.8 < z < 1.9$ (Fig. 1.1), where the cosmic star formation rate history is expected to peak (Cucciati et al., 2012), and where we look forward to a transition between small and disturbed galaxies towards the Hubble sequence. This redshift range is also an opportunity to infer the establishment of the stark dichotomy within the galaxy population, already in place at $z \sim 1$.

MASSIV is an ESO large program (200 hours) grouping 83 galaxies, observed with the near infrared-integral field unit SINFONI on the VLT, from 2008 until completion in 2011 (Contini et al., 2012b). The J- or H-bands have been used to target the redshifted $H\alpha$ line (79 galaxies) and [OIII] line (4 galaxies), both observed with a high spatial resolution ($< 0.8''$) in good seeing conditions, and a total integration time varying between 80 and 120 min. Among these 83 galaxies, 11 were observed with the adaptive optics system (AO), reaching a spatial resolution close to $0.20''$. As of today, it is the largest sample of high- z galaxies observed with IFS in this range of redshift.

The strength of this survey lies in its well defined parent sample, the VVDS (Le Fèvre et al., 2005) a redshift survey selected in magnitude ($I_{AB} \leq 22.5$ for the wide field, $I_{AB} \leq 24$ for the deep field, $I_{AB} \leq 24.75$ for the ultra-deep field) including 35000 spectra in the visible, avoiding any biases linked to a priori color selection techniques. A high completeness of the parent sample is mandatory in order to probe normal and representative galaxies. Galaxies were selected on [OII]3727 equivalent width strength or rest-frame UV intensity from SED fitting, both being a proxy for star formation. This star formation criterion ensures that the brightest rest-frame optical emission line $H\alpha$ ([OIII]5007 for four galaxies) is available to probe resolved kinematics and chemical abundances down to galaxies with a SFR close to $1 M_{\odot} yr^{-1}$. Galaxies were selected also to have a sufficiently close bright star valuable for AO laser guided system.

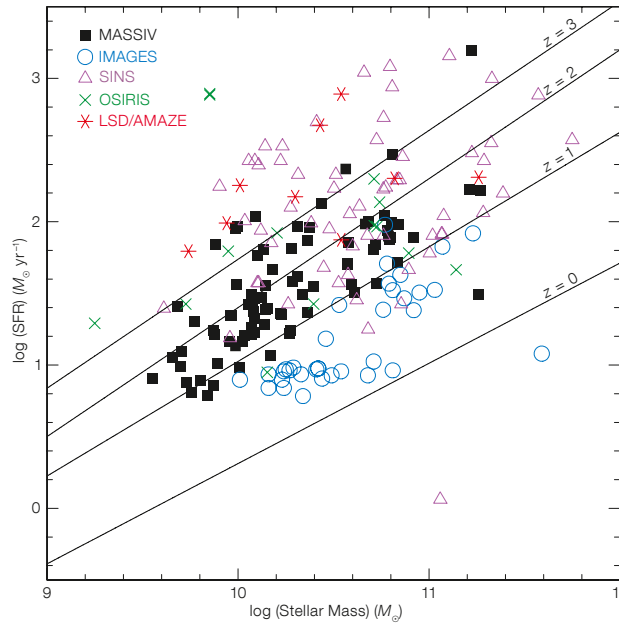


Figure 1.2 – SED-derived star formation rate as a function of stellar mass. The lines represent the empirical relations between SFR and stellar mass for different redshifts between $z = 0$ and $z = 3$ following the analytical expression given in Bouché et al. (2010). All the illustrated IFU samples are rescaled to match the MASSIV sample redshift range. Credits : Contini et al. (2012a).

The continuum I-band magnitude of each galaxy is estimated using the best-seeing CFHT-LS images with a resolution better than $0.65''$.

Figure 1.2 displays the relation between stellar mass and star formation rate for the major IFU samples, over-plotting empirical relations for different redshifts (Bouché et al., 2010). On the one hand, for a given stellar mass SINS, LSD/AMAZE are globally probing galaxies with higher star formation rate than MASSIV. Nevertheless, the dispersion around an average growth of the SFR versus the stellar mass is almost two times lower for the MASSIV sample than for the SINS sample, emphasizing the representativeness of the MASSIV sample. On the other hand, the IMAGES sample is globally probing galaxies with lower SFR at lower redshift for a given stellar mass, as expected. Nevertheless, the stellar mass distribution displays, roughly speaking, two regime of SFR ($\log(\text{SFR}) \sim 1$ & 1.5), unexpected for a representative sample for which we await a continuity in the SFR growth with respect to the stellar mass.

Vergani et al. (2012) investigated fundamental relations using MASSIV data. They derived a gas mass using dynamical arguments, which is estimated on average a fraction of $\sim 45\%$ of the dynamical mass, assuming no central contribution of dark matter (as in Gnerucci et al. (2011)). This choice is consistent with lower concentration parameter of haloes at high- z (Bullock et al., 2001) and is also consistent with the fraction derived using the Kennicutt-Schmidt formulation.

The evolution of the Tully-Fischer Relation (TFR) is expected to be related both to the conversion of gas into stars and to the inside-out growth of dark matter halo by accretion. In Vergani et al. (2012) the stellar mass TFR shows a negligible, net evolution in the past 8 Gyrs with a large scatter that is reduced, but still remarkable, using the S_{05} index ($S_{05} = \sqrt{0.5 \times v_{rot}^2 + \sigma_0^2}$). They interpret this behavior as an evidence of complex physical mechanisms

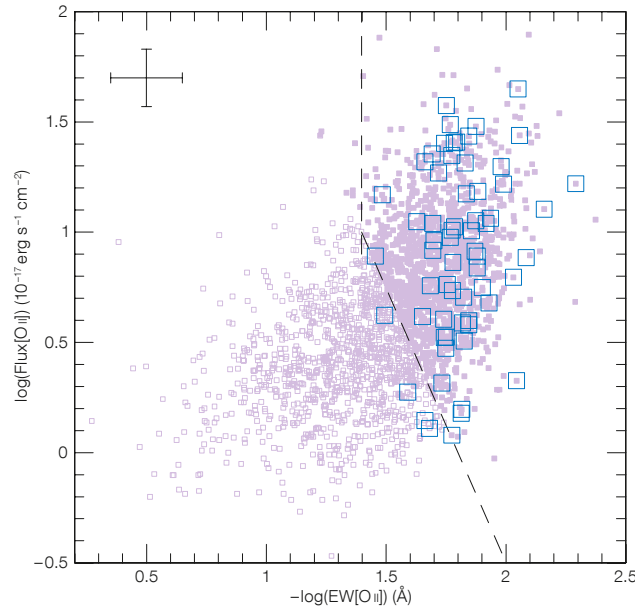


Figure 1.3 – Selection of the MASSIV star-forming galaxies in the WDS parent sample (small grey symbols). The dashed line indicates the selection box based on the [OII] 3727Å emission-line equivalent width, used as a proxy for star formation. This dashed line shows two slopes to account for fainter galaxies that can only be detected with the same signal-to-noise ratio than brighter ones if their continuum decreases as the [OII] flux decreases. The 63 MASSIV galaxies for which the equivalent width is available in the redshift range $z \sim 0.9-1.5$ selected for SINFONI observations are displayed with large blue squares. Credits : Contini et al. (2012a)

at work in the stellar mass/luminosity regime for galaxies in the MASSIV redshift range. They also suggest a marginal evolution in the size – stellar mass and size – velocity relations for which disks become evenly smaller with cosmic time at fixed stellar mass or velocity, and are less massive at a given velocity than in the local Universe, in agreement with cosmological hydrodynamical simulations, e.g. Portinari et Sommer-Larsen (2007).

The first epoch MASSIV sample (50 galaxies) enabled an abundance analysis in Queyrel et al. (2012). The metallicity of galaxies was derived from the [NII]/H α ratio, using the Pérez-Montero et Contini (2009) calibration. Among these 50 galaxies, 26 metallicity gradients were measured inside spatially distinct annular regions defined by the H α contour map. While half of the sample (14 galaxies) is compatible with a null metallicity gradient, a quarter (seven galaxies) of the restricted sample exhibits positive metallicity gradients. Among these positive gradients, four are classified as interacting systems, one might be a chain galaxy, and two are flagged as isolated. Such features have been observed in local interacting galaxies (Werk et al., 2010) and could be due to a fresh metal-poor gas infall onto the core of the merger remnant.

The isolated galaxies detected with positive gradients are more puzzling objects. Such systems have been observed in Cresci et al. (2009) who suggested that cold gas accretion onto the center of the central regions would be a plausible scenario. The radial abundances of the second epoch sample will be published soon (Divoy et al.), and they will broaden up the statistics of this first analysis.

1.2 Kinematical analysis

1.2.1 First epoch sample classification

The full MASSIV sample enables resolved velocity measurement for $\sim 90\%$ of the galaxies. Each velocity field was fitted using ‘flat’ model rotation curve model constructed with a Gaussian kernel mimicking the instrumental Point Spread Function (PSF) (Epinat et al., 2010). A kinematical classification of the first epoch sample has been performed in Epinat et al. (2012), using three distinct estimators. The first parameter which has been used to distinguish fast rotators from slow rotators is the total velocity shear V_{shear} measured on the velocity field, without any inclination correction (see Fig. 1.5). With this simple parameter applied to the nearly finalized sample consisting of 72 galaxies (11 galaxies are undetected), 30 ($\sim 42\%$) galaxies exhibit high velocity shear ($V_{shear} > 100 \text{ km.s}^{-1}$), and 42 ($\sim 58\%$) galaxies have a low velocity shear ($V_{shear} < 100 \text{ km.s}^{-1}$, Epinat et al., to be submitted).

We have also compared a parameter measuring the discrepancy between the position angle of the major axis (PA) of the stellar component and the kinematical PA of the gaseous component with the mean weighted velocity field residuals normalized by the velocity shear (fig. 1.4). Similarly it gives 30 (42%) galaxies classified as “rotating”, and 42 (58%) galaxies classified as “non-rotating” systems.

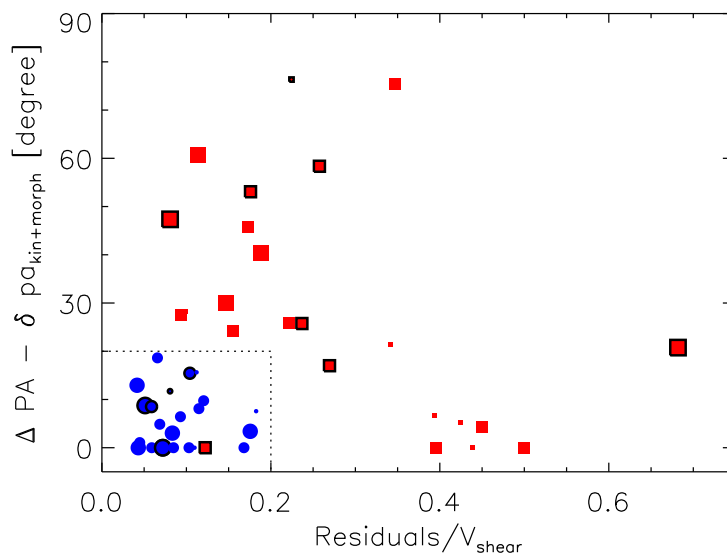


Figure 1.4 – Disagreement between morphological and kinematic position angles as a function of the velocity field residuals normalized by the velocity shear for the galaxies of the MASSIV first epoch sample. Blue dots and red squares represent rotating and non-rotating galaxies. Symbols with black contours mark interacting galaxies. The symbol size is related to the median SNR of the SINFONI data : small, medium and large symbols refer to galaxies with $\text{SNR} < 5$, $5 < \text{SNR} < 10$ and $\text{SNR} > 10$. Credits : Epinat et al. (2012)

In addition, the immediate environment of each galaxy is probed to determine whether the galaxy lies isolated or not. For the entire sample (including undetected galaxies with SINFONI), we have searched for companions in both the H α and I-band images, with a relative velocity lower than 1000 km.s^{-1} and a projected distance lower than $50 \text{ h}^{-1} \text{ kpc}$. Applying this method, 59 (76%) galaxies are found to be isolated, and 19 (24%) are not.

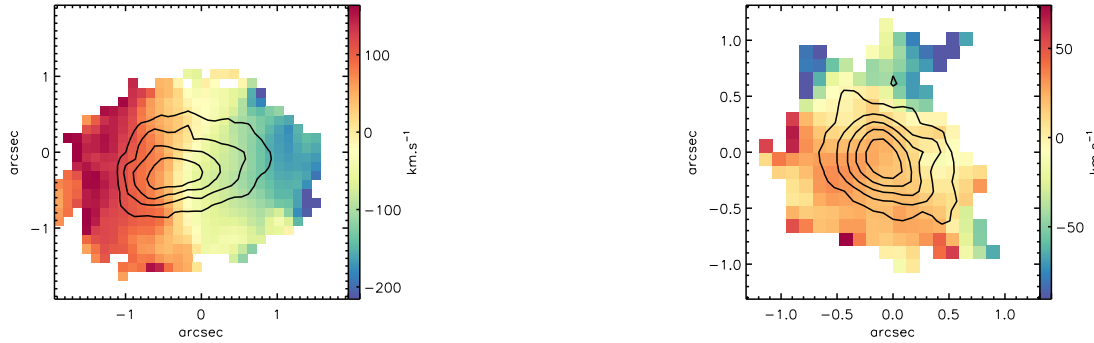


Figure 1.5 – Left : High-shear velocity field (VVDS220584167). Right : Low-shear velocity field (VVDS020386743). Contours are drawing the H α flux distribution. Adapted from Epinat et al. (2012).

This work on close environment has been pushed further in López-Sanjuan et al. (2013). Defining a close pair as a couple of galaxies with a projected radial separation lower than $20 h^{-1}.kpc$, and a radial velocity difference lower than $500 km.s^{-1}$, we have been able to recover a major merger rate $R_{MM} \propto (1+z)^{3.95}$ in the MASSIV redshift range.

Although the origin of galaxies with high V_{shear} is clear (disks with ordered rotation), galaxies with low V_{shear} are more difficult to understand. The possible explanations could be : face-on disks, unrelaxed merger remnants, ionised gas distribution dominated by large star-forming clumps, or star-forming spheroids. Nevertheless, the orientation of disks with spin vectors randomly distributed could only account for 14% of this low-shear population considering disks rotating at $V_{rot}=200 km.s^{-1}$. The large fractions of interacting galaxies, and of non-rotating galaxies seem to suggest that mass redistribution mechanisms are at work in the redshift range $1 < z < 2$.

1.2.2 Kinemetry

This thesis aimed at developing a new method of analysis based on an harmonic decomposition technique that could provide a new kinematical classification. Indeed, harmonic decomposition techniques provide different terms that disentangle the rotation from non circular motions. In addition, different kinds of non circular motions nature could be identified. Such an analysis has been performed by Shapiro et al. (2008) on the SINS sample. I used this work as a starting point for the new analysis of the MASSIV sample that relies on the kinemetry method (Krajnović et al., 2006). The kinemetry algorithm is based on the observation that the projected circular velocity at any location in the plane of a disk galaxy can be written as

$$V_{obs}(r, \theta) = V_{sys} + V_{rot}(r) \cos(\phi) \sin(i), \quad (1.1)$$

where r and ϕ are the polar coordinates in the galactic plane, i is the disk inclination with respect to the line-of-sight. The kinemetry is an algorithm that consists in performing an harmonic expansion along ellipses drawn on the velocity map around the kinematical center. This information can be easily modeled with a decomposition by a finite number of harmonic terms :

$$V_{kin}(r, \phi) = A_0(r) + \sum_n \left[A_n(r) \sin(n\phi) + B_n(r) \cos(n\phi) \right], \quad (1.2)$$

where V_{kin} is the reconstructed radial velocity, r is the value of the semi-major axis of the ellipse, ϕ is the azimuthal angle in the plane of the galaxy, A_n and B_n are the amplitude terms of the n -th harmonic.

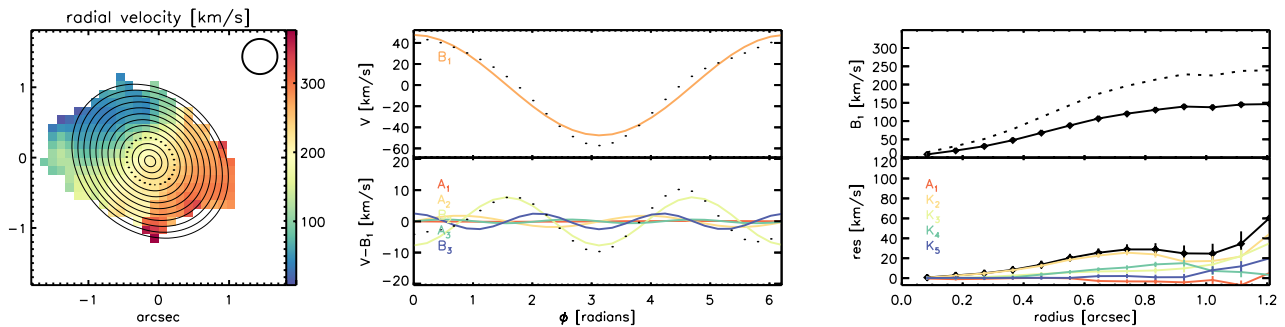


Figure 1.6 – Illustration of a kinemetry run on the velocity field of VVDS220376206, a galaxy from the MASSIV sample. On the left : Velocity field of the toy disk model, with kinemetry ellipses over-plotted. The circle on the top right-hand corner represents the size of the PSF (FWHM). In the middle : Kinemetry harmonic expansion as a function of the azimuthal angle along the ellipse with black dots on the velocity field. The top panel shows the radial velocity measurements (black dots), and the fit to the kinemetry rotation velocity term (B_1 , orange line), while the lower panel shows the velocity residuals once the rotation has been subtracted ($V - B_1$, black dots) and the fit to the higher order harmonic coefficients (A_1, A_2, B_2, A_3, B_3). We clearly see on this plot that the term B_1 has the highest velocity amplitude. On the right : Kinemetry expansion for all the ellipses. The top panel shows the radial dependence of the velocity rotation term (black solid line) and the deprojected velocity rotation (black dotted line), which correspond to the classical rotation curve. The lower panel shows the radial dependence of the residual coefficients A_1 and K_n , $n=2$ to 5 (see 1.5). The black line shows the quadratic sum of all the residual coefficients up to the order $n=5$. The stronger residual is seen in coefficient K_2 which contains B_2 .

Equation 1.1 can easily be modeled by equation 1.2 since they both have a $\cos \phi$ term, and both are expressed in the same polar coordinate system. Therefore, the algorithm consists in finding the best ellipses parameters $[PA, i, x_c, y_c]$ which minimize the amplitude of the harmonic terms that does not correspond to a circular motion. The best-sampling ellipses are found by minimizing the quantity which is restricted to the first three orders of the harmonic expansion.

The kinemetry algorithm is able to determine a best-sampling ellipse, for a given semi-major axis defined by a PA and an axis ratio

$$q = \frac{b}{a} = \cos i \quad (1.3)$$

with an inclination i , that minimizes the quantity

$$\chi^2 = A_1^2 + A_2^2 + B_2^2 + A_3^2 + B_3^2, \quad (1.4)$$

The variation of PA and q with radius is justified for kinematical data with a very high signal-to-noise ratio (SNR) and is typically like the one encountered in the galaxies of the nearby

Universe. Especially, it is interesting to use this adjustment on HI data that often display disk warps in the neutral component. Adapting the geometrical parameters of the ellipses at large scale allows us to extract the underlying rotation in a more robust way. However, the poor resolution of the high redshift IFU data only allows to sample the optical disk and thus requires to reduce the number of free parameters. Each kinemetry expansion in this work is therefore performed with all the geometrical parameters kept constant, just as it has been done in Shapiro et al. (2008). Using in such a way the kinemetry method enables to quantify the deviation from the case of a purely rotating disk by estimating the power in the harmonic coefficients other than B_1 . Krajinović et al. (2006) define the coefficients k_n which are the quadratical sum of the even and odd coefficients of a given order n of the harmonic expansion :

$$k_n = \sqrt{A_n^2 + B_n^2} \quad (1.5)$$

in order to quantify the residual coefficients of a given harmonic n . An illustration of a kinemetry expansion technique on the velocity field of a galaxy from the MASSIV sample is presented in figure 1.6.

1.2.3 Point Spread Function correction

Measuring the disk inclination on such low spatial resolution data is critical for this work. A good inclination estimate is needed to recover properly the terminal rotational velocity V_c^{max} , that is expected to study fundamental relations and their evolution through cosmic times (e.g. the Tully-Fischer relation). Drawing ellipses on a velocity field with a constant axis ratio to extract the kinemetry coefficients is not a satisfying approach when taking into account the poor resolution. Indeed, the PSF being quite extended regarding the galaxy size in the SINFONI high-redshift observations, the PSF effect on the shape of the ellipses superimposed to a velocity map should be quantified.

To estimate the PSF impact on a velocity field, I have produced multiple toy models of an thin exponential disk, which is a good first order approximation to model the central galactic regions of a rotating disk. The radial surface density of such disk is written

$$\Sigma(r) = \Sigma_0 \exp\left(-\frac{r}{r_d}\right) = \frac{M_d}{2\pi r_d^2} \exp\left(-\frac{r}{r_d}\right), \quad (1.6)$$

with Σ_0 the central surface density, M_d the total disk mass, and r_d the disk scalelength. The circular velocity for an exponential disk in a dynamical equilibrium state is an analytical function, and is written

$$v_c(r) = 4\pi G \Sigma_0 \frac{r^2}{4r_d} \left[\mathbb{I}_0\left(\frac{r}{2r_d}\right) \mathbb{K}_0\left(\frac{r}{2r_d}\right) - \mathbb{I}_1\left(\frac{r}{2r_d}\right) \mathbb{K}_1\left(\frac{r}{2r_d}\right) \right], \quad (1.7)$$

with \mathbb{I}_0 , \mathbb{K}_0 , \mathbb{I}_1 , \mathbb{K}_1 being respectively the modified Bessel functions of the order one and two, and G the gravitational constant (Freeman, 1970). Using equations 1.6 and 1.7, I am able to build 2D surface density and radial velocity maps for a given set of geometrical parameters (PA, i, x_c, y_c) positioning the disk in the sky plane. The produced maps adopt the pixel size of the SINFONI with AO mode observations (0.05"). I have constructed the velocity maps of these disk models using a Gaussian kernel mimicking the instrumental PSF. Following Epinat et al. (2010), the convolved velocity field is built in modulating the raw velocity model by the flux distribution and the spatial PSF using the following relation :

$$V_{conv} = \frac{[V \times F] \otimes_{xy} \text{PSF}_{xy}}{F \otimes_{xy} \text{PSF}_{xy}}, \quad (1.8)$$

where V_{conv} is the convolved 2D velocity map, V is the 2D velocity map, F is the 2D H α flux map, PSF_{xy} is the 2D Gaussian kernel modeling the PSF and where \otimes_{xy} is the 2D convolution operator. The H α monochromatic flux map is estimated to be proportional to the surface density map, and the flux is scaled to match Bouché et al. (2010) SFR-mass relation for a given redshift. The convolution is performed in the Fourier plane, where the operation consists in a simple product :

$$F \otimes_{xy} \text{PSF}_{xy} = \mathcal{F}^{-1} \left(\mathcal{F}(F) \times \mathcal{F}(\text{PSF}_{xy}) \right) \quad (1.9)$$

where \mathcal{F} and \mathcal{F}^{-1} are the direct and inverse Fourier transform functions respectively. To avoid low sampling issues during the convolution operation, I insure that the flux and radial velocity map are sampled by pixels four times smaller than the PSF full width at half maximum (FWHM). The convolved maps are resampled after the convolution, in order to obtain the required pixel size. In order to explore the spatial resolution dependance, a set of disk toy models have been constructed with the following parameters :

- 7 disk scalelength with $r_d=[2,3,4,5,6,7,8]$ kpc
- 12 PSF FWHM with $s=[0.1,0.2,0.3,0.4,0.5,0.6,0.7,0.8,0.9,1.0,1.1,1.2]$ arcsecond
- 3 disk inclination with $i=[40,60,80]$ degrees.

The resulting $7 \times 12 \times 3 = 252$ exponential disk toy models are presented in figure 1.7.

In a first step, I ran a kinemetry expansion with a constant ellipse flattening corresponding to the underlying disk inclination i . Such an expansion is displayed in figure 1.8 (top line). We observe that the power in the B_3 kinemetry term is significant. Thanks to Krajnović et al. (2006), we know that coefficients $B_{2n+1, n>1}$, hence B_3 , can be interpreted as the direct signature of the PSF convolution with the pure rotating disk velocity field. Hence, it is possible to compensate the beam smearing effect by adjusting at each radius the flattening of the kinemetry ellipse as illustrated in figure 1.8 (bottom line). Indeed, isophotal axis ratio is always underestimated in ground-based morphology observations without AO of high redshift galaxies due to the seeing FWHM s , which is only a couple of time smaller than the galactic radius. In addition, external isophotes have a lower axis ratio than central ones. This led Epinat et al. (2010) to define the axis ratio affected by the convolution q_c from the one of the underlying disk q given in equation (1.3) in introducing a coefficient c to model the fraction of the seeing that should be added on to the actual major and minor axis to account for seeing effect. In order to isolate the axis ratio taking into account the seeing effect rather than the actual axis ratio, I have inverted the equation (7) in Epinat et al. (2010) :

$$q_c = \frac{b_c}{a_c} = \sqrt{\frac{b^2 + c^2 \times s^2}{a^2 + c^2 \times s^2}}. \quad (1.10)$$

with a and b are respectively the actual major and the minor axis of the underlying disk ; a_c and b_c being respectively the major and minor axes of the ellipse including seeing effect. In expression (1.10), a and b are both quadratically overestimated by a fraction c of the seeing s and $q_c(r)$ is the flattening of an ellipse which minimizes the residuals in the harmonic expansion.

The coefficient c could be isolated from equation (1.10), introducing $r = \frac{a_c}{2}$, the semi-major axis ratio of the kinemetry ellipse, we have :

$$c = \frac{2r}{s} \sqrt{\frac{q^2 - q_c^2}{q^2 - 1}}, \quad (1.11)$$

In the kinemetry expansion, I have used this prescription instead of using a constant flattening for the kinemetry ellipses. Therefore, the axis ratio of an ellipse drawn on the PSF-convolved velocity field, which minimizes the $B_{2n+1, n>1}$ coefficients, is described by relation (1.10). Assuming that the exponential disk is a good first order approximation for a high redshift rotating disk, I seek a function for c which could minimize the B_3 terms in the kinemetry expansion. I modified the original kinemetry routine of Krajnović et al. (2006) in order to only keep the kinemetry ellipse flattening q_c free during the brute-force minimization of equation 1.4, hence of the B_3 term, as illustrated in figure 1.8 (bottom line). Then I ran the kinemetry algorithm on each of the toy models presented in figure 1.7 and I recovered the value of c for all the semi-major axis values r of the ellipses drawn by the kinemetry.

The function c is expected to depend on four parameters : the inclination and the scalelength r_d of the underlying disk, the seeing s and the semi-major axis of the considered kinemetry ellipse r . The recovered values of c exhibit an ordered pattern across many different toy model parameters, without a clear dependence with inclination. In addition, the c functional form depends on dimensionless quantities because this problem is fully scalable. Therefore, I have empirically searched for a function that fits the measurements of c and that depends on the dimensionless ratios s/r_d and r/r_d . I have visually adjusted the following best analytical parametrization :

$$c\left(\frac{r}{r_d}, \frac{s}{r_d}\right) = \left(\frac{3}{2} + \frac{r}{5r_d}\right) \left[1 - \exp\left(-\frac{r}{r_d} \left[\frac{r_d}{4s} + 1\right]\right)\right], \quad (1.12)$$

The measurements of c are displayed in figure 1.9 (top) as well as the best parametrization model of c (bottom). This empirical analytic function reflects the convolution effect on the fraction of the seeing that should be added to the major and minor axis of a given ellipse. With this formulation, the residuals due to bad flattening never exceed 5% of $V_c(r)$ on all the presented toy models. We note that this equation is governed by the value of the seeing, but also by the value of the disc scalelength. The correction of the PSF effect on the shape of the kinemetry ellipses thus requires an estimate of the disk scalelength that can be provided by morphological analysis. If this measurement is not available, the rotation curve can first be extracted from the velocity field, then an exponential disk rotation curve can be fitted, and finally get an estimate of the disk scalelength in order to be able to run the kinemetry algorithm with ellipses flattening corrected from the PSF.

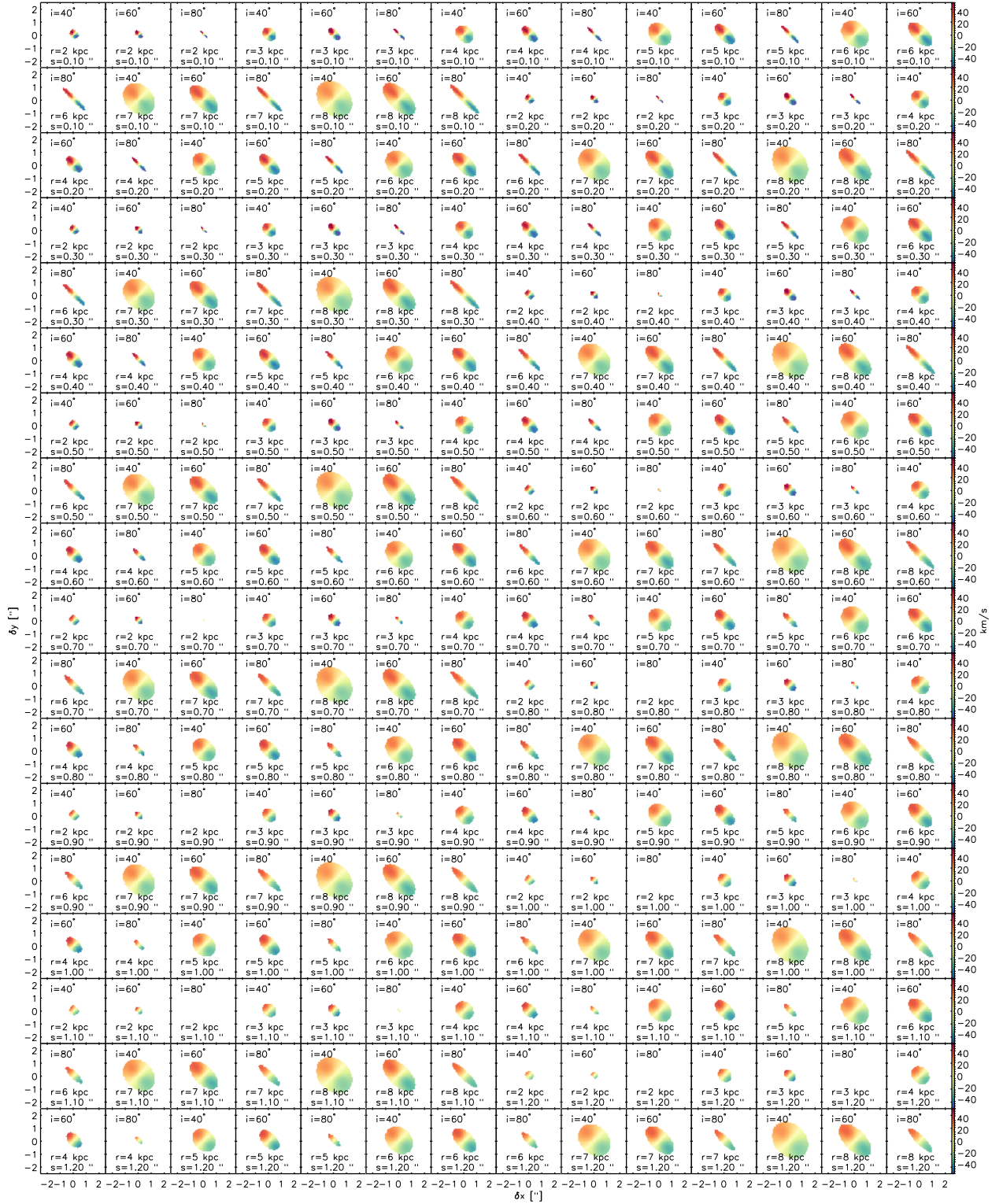


Figure 1.7 – Velocity fields for exponential disk toy models. The set of 252 toy models is exploring 3 disk inclinations $i=[40,60,80]$ degrees, 7 disk scalelength $r_d=[2-8]$ kpc and 12 PSF FWHM $s=[0.1-1.2]$. The PA is equal to 45° , the pixel size is equal to $0.05''$.

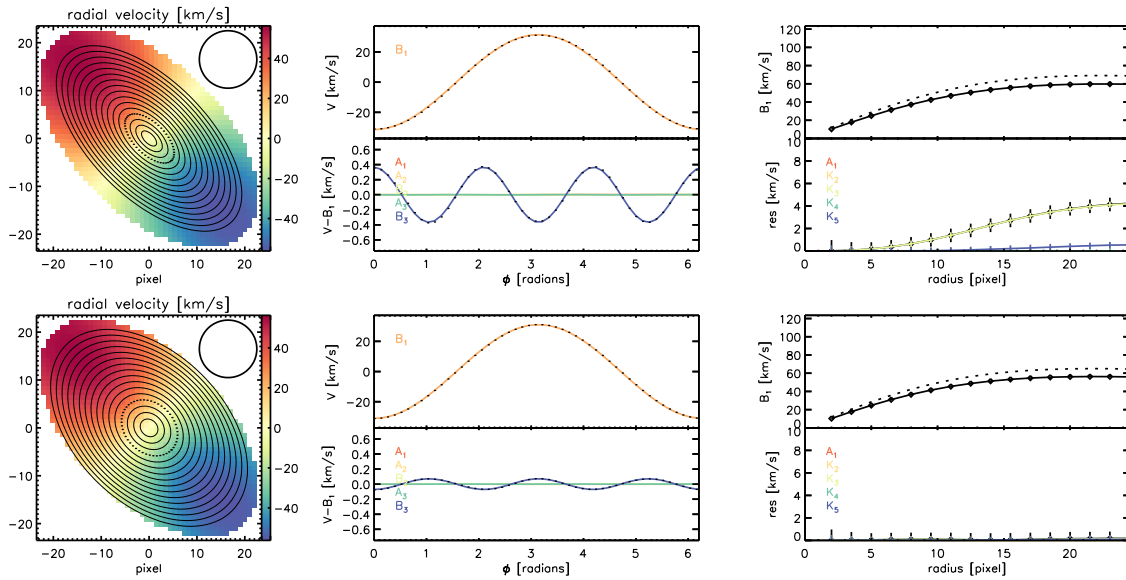


Figure 1.8 – Same as figure 1.6 but in order to illustrate the PSF effect on a toy model exponential disk velocity field with an inclination $i=60^\circ$, a scalelength $r=3$ kpc and a PSF FWHM $s=0.60$ arcsec. On the left : Velocity field of the toy disk model, with kinematics ellipses overlaid. In the middle : Kinematic harmonic expansion as a function of the azimuthal angle. On the right : Kinematic expansion as a function of the radius (ellipses). The PSF FWHM is represented by black circles on the top right-hand corner of each velocity field. The bottom line displays the case without PSF correction for which all the ellipses have the same axis ratio. The top row shows the case with PSF correction where the ellipse axis ratio increases from the center to the outskirts; combining the PSF and the disk scalelength, each ellipse flattening has been adapted to minimize all the harmonic coefficients except B_1 .

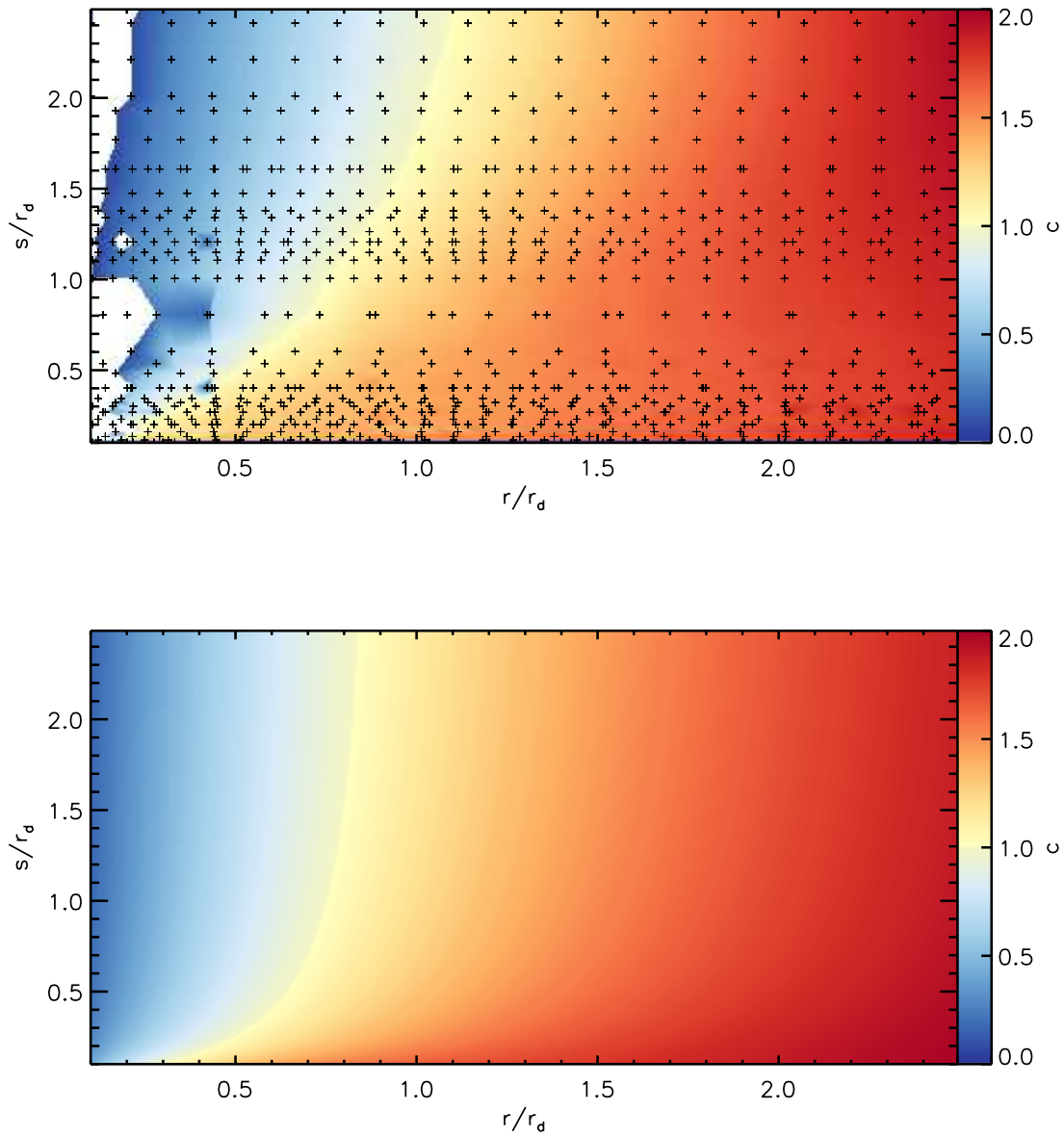


Figure 1.9 – PSF contribution to the ellipse minor axis function $c(\frac{r}{r_d}, \frac{s}{r_d})$ determined empirically with the data set of exponential disk toy model. At the top : measurements of c on the 252 toy models. Black crosses represent the sampling of the parameter space with the toy models and the underlying colormap displays the values of c triangulated on a regular high-resolution grid. At the bottom : best analytical parametrization model of c .

1.2.4 Geometrical parameters fit

In order to estimate the disturbance of a given velocity field, we need a robust velocity field modeling algorithm able to adjust the geometrical parameters $[PA, i, x_c, y_c]$. The X^2 function of a velocity field model is characterized by multiple local minima. Gradient-based non-linear X^2 minimization methods (e.g. Levenberg-Marquardt algorithm) are very sensitive to the initial values of the model parameters, and can be quickly trapped into these local minima of the X^2 function. This induces a very poor quality of the geometrical parameters adjustment. In order to avoid these spurious effects in the minimization, implied by the degeneracy of the modeling, we need an algorithm which explores the parameters space in the most efficient and possible way. Those constraints lead us to adopt a genetic algorithm, which fulfills all the requirements for this kind of analysis. Although the choice of the minimization algorithm is inspired from the work of Cresci et al. (2009), our method differs in the technique used to model the velocity fields. While Cresci et al. (2009) are using an exponential disk model, we use the kinemetry rotation harmonic coefficient B_1 to produce a model velocity field. This approach has the advantage of assuming no rotation curve model, and we expect better results on the quality of the recovery of geometrical parameters, especially in the case of a rotating disk with radial motions in the disk plane, which could lead the exponential disk modeling towards a wrong PA estimation.

We wrapped the *Pikaia* genetic algorithm implementation (Charbonneau, 1995) into an IDL analysis code named *KINEGEN* in order to obtain a strong heuristic optimization method. *KINEGEN* is one of the products of this thesis, and which have been made available to the scientific community¹. The documentation of this code is presented in the appendice C. The *Pikaia* genetic optimization code works with a population of individuals (i.e. phenotype), each of them being characterized by a gene (i.e. genotype), that encodes the values of the model parameters. Each individual is characterized by a fitness function, that evaluates the model residuals. As a matter of fact, the fitness function is proportional to the inverse of a X^2 function. The genetic algorithm maximizes the fitness function (i.e. minimizes the X^2 function) of a population of phenotypes by performing the following recursive steps :

1. Build a random initial population composed of n_{pop} individuals and evaluate the fitness f of its members. The number of individuals remains constant throughout the entire process, so that individuals with a low fitness disappear from the initial population.
2. Generate a child population by "breeding" selected pairs of individuals from the old population. The probability of an individual to be selected for breeding is based on its fitness : individuals with a high fitness are more frequently selected for breeding than individuals with a low fitness. As theorized in the Theory of Evolution, the breeding process is composed of two distinct genetic events : the crossover of two genes from two parent individuals with a probability p_c , and the spontaneous mutation of a gene with a probability p_m .
3. Evaluate the fitness of each individual of the new population.
4. Insert the new offspring into the old population, and remove the lowest fitness individuals to keep n_{pop} constant.
5. Check convergence ; unless fittest phenotype matches target phenotype within tolerance, go back to step (2). The recursion is stopped if the generation number exceed a threshold n_{gen} .

1. <http://code.google.com/p/kinegen/>

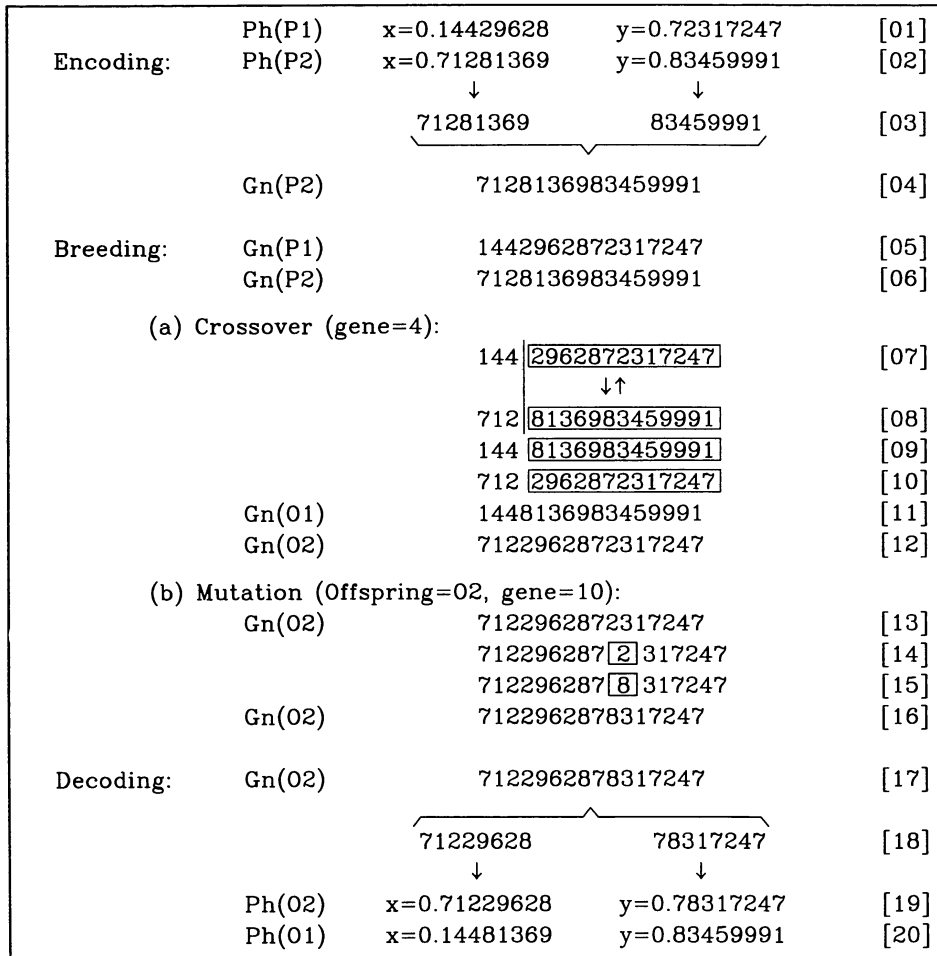


Figure 1.10 – Encoding, breeding and decoding in genetic algorithm. In this example, phenotypes are defined in terms of two parameters x and y (real numbers) of a given model and are encoded as a string of 16 decimal digits. "Ph(P1)" means phenotype of parent P1, "Gn(O2)" is "genotype of offspring 2", and so on. Encoding is only shown for Ph(P2), and decoding for Gn(O2). In the presented implementation, a breeding event produces two offspring, and both crossover and mutation occur only if a probability test yields true. Credits : Charbonneau (1995)

Applying this genetic scheme on a population of individuals forces the genotypes to converge towards a state of best fitness population through selection pressure. The selection pressure is controlled by the crossovers rate, the mutations rate and the breeding occurrences. The breeding process is illustrated in figure 1.10. In the case of the modeling of a velocity field, the fitness function for a given set of geometrical parameters should be proportional to

$$f(p) \propto (n_{pix,kin}(p) - n_{params}) \sum_i \left[\frac{\sigma_i}{V_{kin,i}(p) - V_i} \right]^2, \quad (1.13)$$

where p is the set of geometrical parameters set, V_i is the i -th velocity measurement, $V_{kin,i}$ is the i -th associated modeled velocity, $\sigma_{kin,i}$ the i -th associated errors to the measurement of

emission line positions, $n_{pix,kin}$ is the number of pixels recovered by the kinemetry expansion, and n_{params} is the number of free parameters.

In the case of the KINEGEN algorithm, the modeled velocity is obtained by performing a kinemetry expansion on the input velocity field for a given set of geometrical parameters p . The kinemetry method does not necessarily model all the pixels of the input velocity field because it uses ellipses to extract the kinematical information and these ellipses do not necessarily cover the full input velocity field. Thus it is required that the fitness function rewards genotypes where the kinemetry expansion covers a high number of pixels. Strong offsets between kinematical and morphological centers are not expected. In order to avoid such a situation, the fitness function should also reward genotypes that provide kinematical center close to the peak of the emission line flux. Therefore, we define a merit factor \mathcal{M} by the expression

$$\mathcal{M} = \left(\frac{1}{r_{peak}(p)} \right)^{\gamma_1} \left(\frac{1}{n_{pix,kin}(p)} \right)^{\gamma_2} \quad (1.14)$$

where r_{peak} is the distance of the input center to the peak of H α flux and where γ_1 and γ_2 are reward coefficients. This merit factor is incorporated in the fitness function evaluation in order to boost the fitness of solutions, that fulfill the reward criteria :

$$f(p) = \mathcal{M} \times (n_{pix,kin}(p) - n_{params}) \sum_i \left[\frac{\sigma_i}{V_{kin,i}(p) - V_i} \right]^2, \quad (1.15)$$

The solutions that are found close to the H α flux peak are boosted with the reward coefficient γ_1 . The kinemetry expansions that produce a high number of pixels in the velocity field model are boosted with the reward index γ_2 in the above formula. A high value for γ_1 means a high probability to find a kinematical center close to the position of the maximum H α flux. A high value for γ_2 means a high probability to find geometrical parameters that enable kinemetry ellipses to cover completely the velocity field. In order to prevent the convergence of the genetic algorithm towards a wrong model, the user of the KINEGEN routine can introduce some "fail" conditions which translate into a fitness $f=0$ for the concerned individuals. I have implemented three "fail" conditions in KINEGEN :

- The ratio of the number of pixels modeled over the number of pixels in the input should be greater than a minimal fraction of recovered pixels (\min_{pix}).
- The slope of the linear regression on the radial rotation term B_1 should be greater than a lower limit (\min_{slope}).
- The rotation term deprojected from the recovered inclination should be lower than the maximum deprojected rotational velocity (v_{limit}).

Although these parameters are optional, they can help to accelerate the convergence towards a good value for the geometrical parameters.

1.2.5 Exponential disk toy model test case

In order to test the reliability of the KINEGEN routine described in the previous section, I have produced a set of SINFONI mock observations of an exponential disk toy model with a scalelength $r=3.5$ kpc and a mass $M=10^{11}M_{\odot}$ (corresponding to a maximum circular velocity

of $\sim 220 \text{ km.s}^{-1}$). The procedure to produce such toy models disk is similar to the one described in the section 1.2.3. The outputs of the toy model generation routine are λ -cubes similar to SINFONI observations, with a spatial sampling of $0.125''$, a spectral sampling of 1.95\AA , a spatial resolution of $0.60''$ and a spectral resolution of 45 km.s^{-1} .

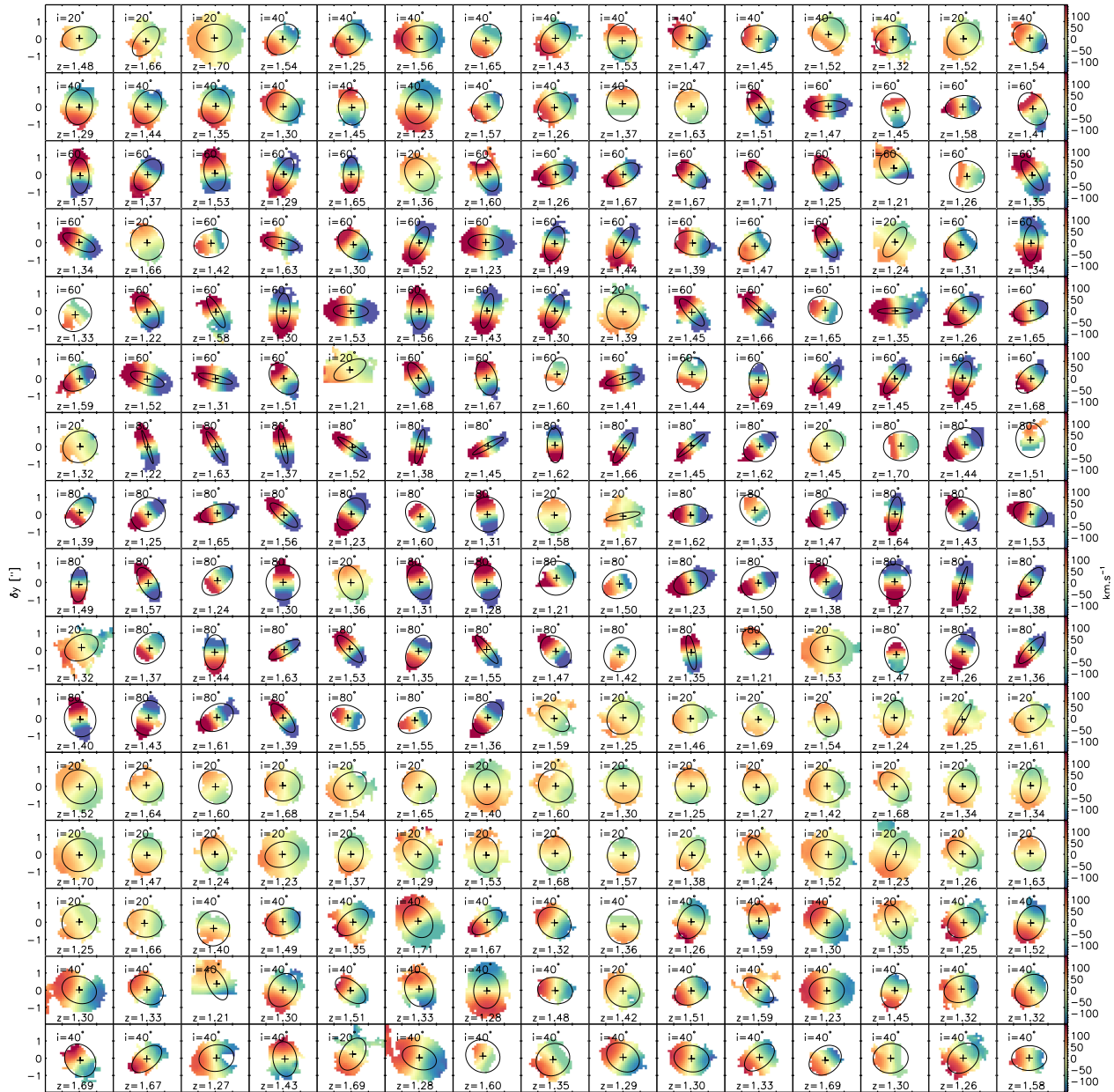


Figure 1.11 – Velocity fields of the high-SNR disk toy models. For each toy model, the redshift and the input inclination are indicated. The result of the geometrical parameters adjustment is also illustrated on each velocity field with a black cross that indicates the recovered kinematic center and with an ellipse that indicates the recovered PA and inclination.

To obtain a realistic power spectrum for the typical noise of a SINFONI observation, I have inserted the toy model λ -cubes into a flux calibrated SINFONI λ -cube from the MASSIV sample, using the VVDS020106882 observation. I model the $H\alpha$ emission line at a random

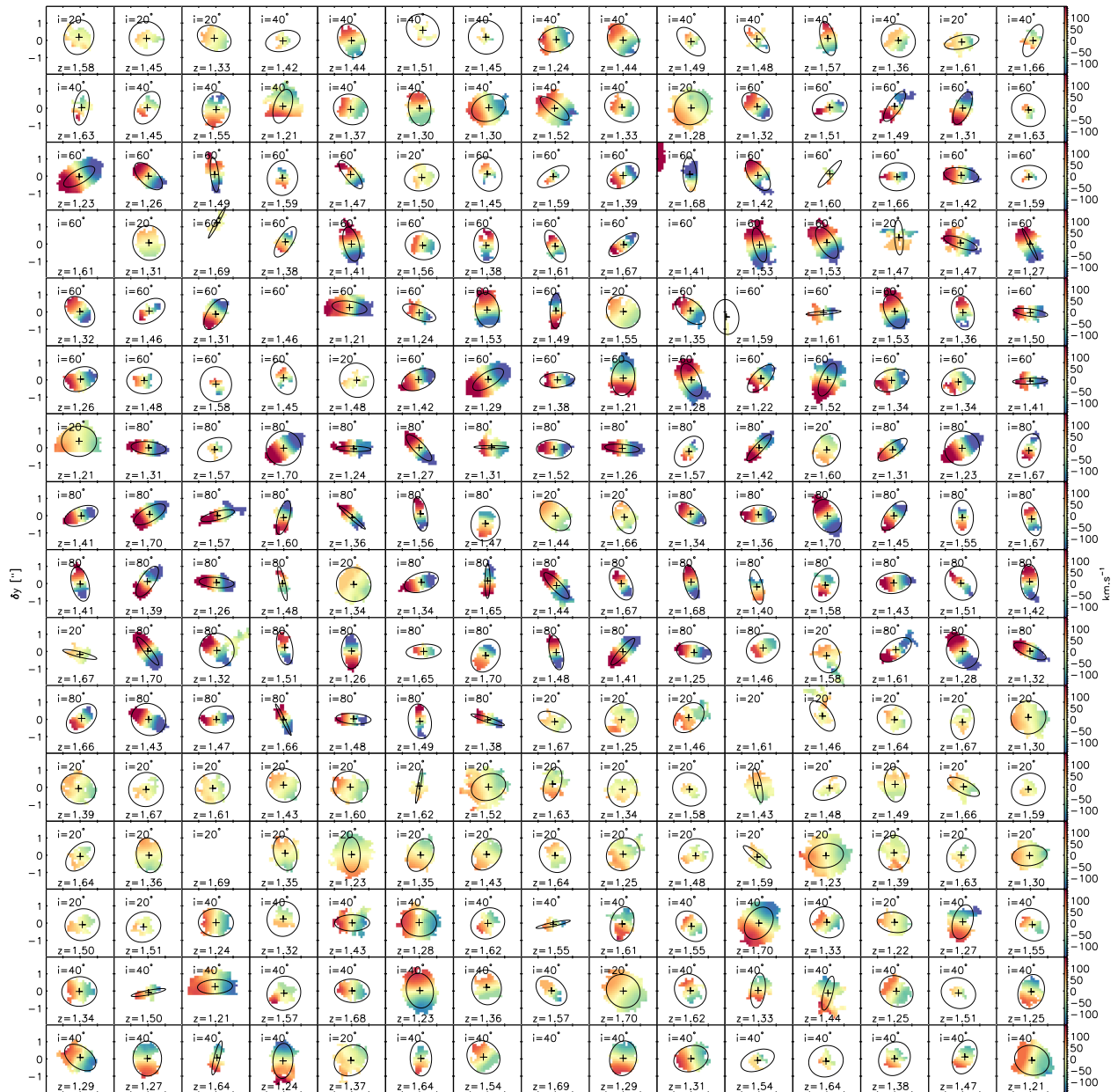


Figure 1.12 – Velocity fields of the low-SNR disk toy models. For each toy model, the redshift and the input inclination are indicated. The result of the geometrical parameters adjustment is also illustrated on each velocity field with a black cross that indicates the recovered kinematical center and with an ellipse that indicates the recovered PA and inclination.

spectral position between redshift 1.21 ($\lambda=1.45\mu\text{m}$) and redshift 1.71 ($\lambda=1.78\mu\text{m}$) to obtain different noise realizations. I taken care to avoid the redshift $z=1.399$ where the real galaxy is located. I have built two distinct sets of toy models, where a first set of high-SNR λ -cubes have their flux scaled to reach a fixed signal-to-noise ratio $\text{SNR}_{\text{max}}\approx 30$ in the pixel that contains the highest flux value. The second toy model low-SNR λ -cubes have their flux scaled to reach $\text{SNR}_{\text{max}}\approx 10$. These two maximum values for the SNR are comparable to the highest and lowest SNR in the MASSIV sample, allowing direct comparisons. The choice of keeping the maximum SNR constant with respect to the inclination, is motivated by the fact that we do not see any SNR inclination dependence in the MASSIV sample (see figure 1.13). Indeed, edge-on disks are more subject to absorption than face-on disks, where the flux should be more extended spatially if the gas is considered to be optically thin.

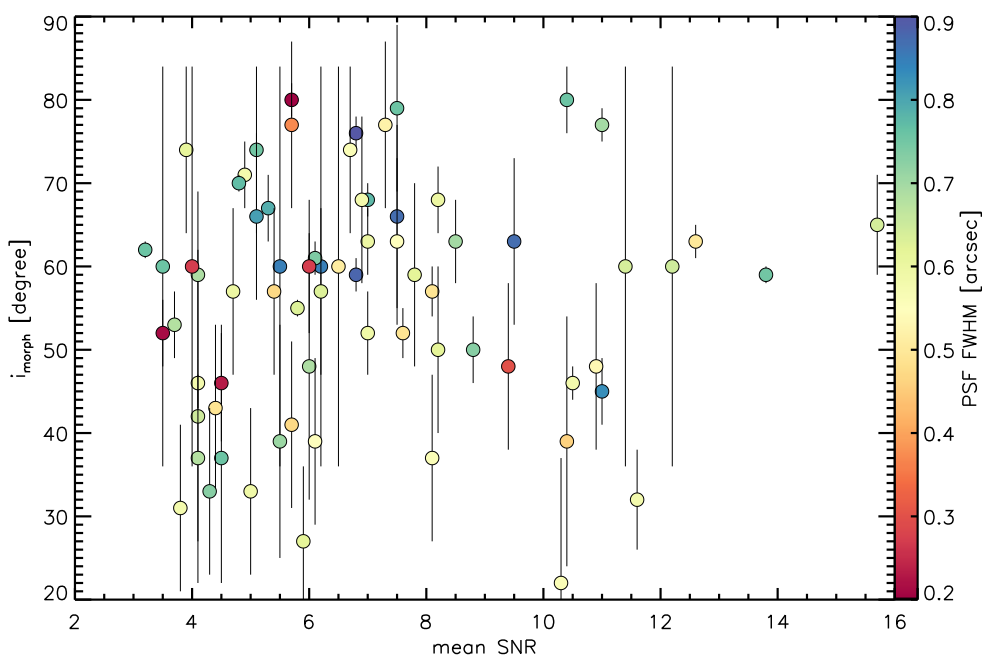


Figure 1.13 – Distribution of disk inclinations in the full MASSIV sample, measured on the stellar morphological data (CFHT I-band images) as a function of mean $\text{H}\alpha$ SNR value in areas with signal detected beyond a confidence level of 3σ . The color of each symbol encodes the PSF size of the SINFONI observation.

Both of the toy models datasets are composed of 240 mock observations. In each toy model, the PA is picked randomly, while the kinematical center is set to the center of the λ -cube. I deliberately choose to not pick randomly the toy model disk inclination because I wanted to test the ability of the KINEGEN algorithm to recover the input disk inclination, that is the most sensible parameter to recover. I have set four distinct values of the inclination $i=[20^\circ, 40^\circ, 60^\circ, 80^\circ]$, each of them being sampled by 60 toy models with random PA and redshift.

Once the toy models are built, the resulting λ -cubes are analyzed with the numerical tools developed for the MASSIV sample. The analysis consists in adjusting a Gaussian emission line ($\text{H}\alpha$ in our case) in each spectrum. The achieved velocity fields are presented in the figures 1.11 and 1.12 respectively displaying the high and low SNR toy models. As we want to run

the KINEGEN routine only on useful pixels, these velocity fields only display the pixels with a SNR greater than 5. However, the presence of some artefacts due to the noise and bad sky lines subtraction in the λ -cubes can make the Gaussian fitting procedure end up with SNR higher than 5 in regions where the disk toy model does not contribute to the H α flux. On a small dataset, these bad pixels can manually be removed one by one, but this method is not possible on large data set, as the one we study here.

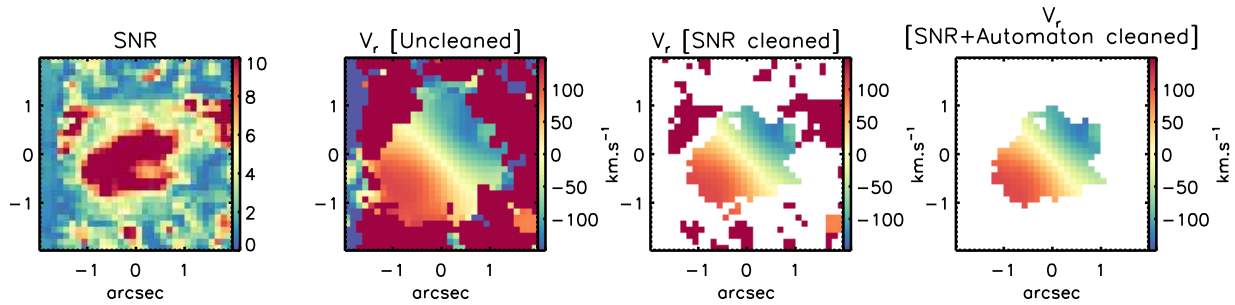


Figure 1.14 – Illustration of the cleaning procedure of moment maps of a high-SNR disk toy model. From left to right : SNR map, uncleaned velocity field, velocity field without pixels with SNR<4, velocity field without pixels with SNR<4 and cleaned by the cellular automaton.

To remove these false signal detections, we used a cellular automaton algorithm which detects spatially connected pixel groups. The cellular automaton wanders from one pixel to another in the SNR cleaned velocity field map, following specific displacement rules. These rules stipulate that the cellular automaton can only explore pixels (i) that are connected to the current pixel explored (ii) that are not explored yet (iii) that encode a velocity value closer than 25 km.s⁻¹ from the current value. With these rules, we are able to detect each interconnected pixel groups. I have removed all the groups that are composed of less than 25 percent the total number of pixels with a velocity value in the SNR cleaned velocity map. An illustration of this two-pass cleaning process is presented in figure 1.14.

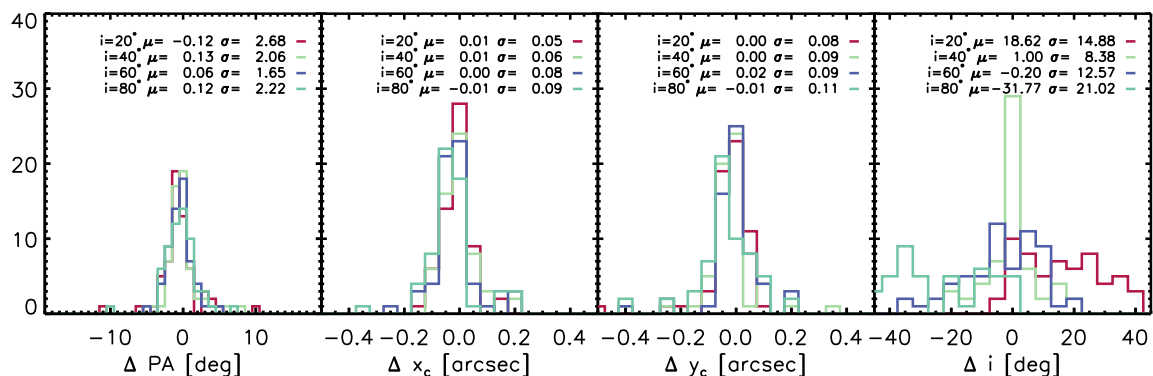


Figure 1.15 – Distribution of the recovered geometrical parameters subtracted from the input values for the high-SNR toy model data set. On each panel, the median value μ and the dispersion σ of the recovered parameters for a given input inclination is indicated.

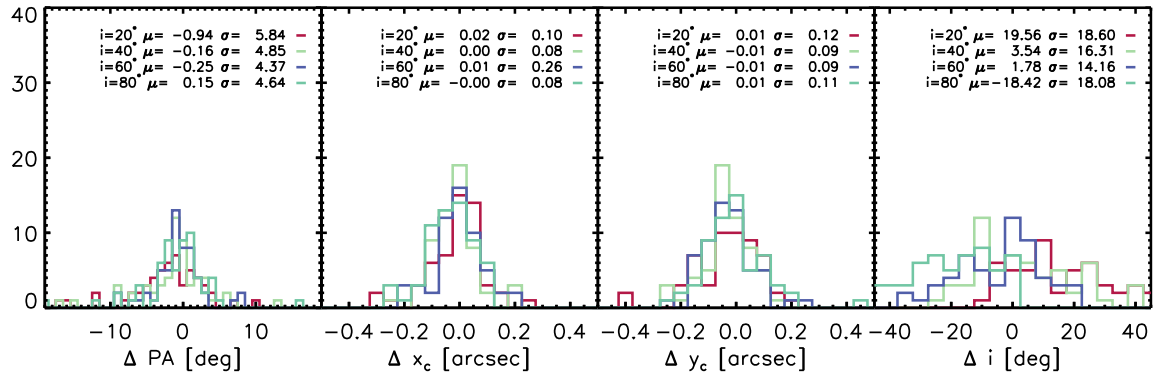


Figure 1.16 – Distribution of the recovered geometrical parameters subtracted from the input values for the low-SNR toy model data set. On each panel, the median value μ and the dispersion σ of the recovered parameters for a given input inclination is indicated.

I have run the KINEGEN algorithm on each toy model velocity field. The resulting recovered geometrical parameters are illustrated with crosses and ellipses in figures 1.11 and 1.12. The following strategy was applied to fit the geometrical parameters :

- In a first-pass fit, I recover all the geometrical parameters by setting the reward coefficients to $\gamma_1 = \gamma_2 = 6$. The genetic algorithm is set to produce $n_{pop} = 150$ individuals over $n_{gen} = 15$ generations with a crossover rate $p_c = 80\%$ and a mutation rate of $p_m = 20\%$, that ensures a good convergence for the geometrical parameters.
- In a second-pass fit, I recover a better value for the inclination by performing the same fit but keeping constant the PA and center to the value found in the first-pass fit. I set $\gamma_1 = 0$ so that the algorithm is no more rewarded for finding solutions with more pixels. I lower the kinemetry cover parameter to 35%, so that each kinemetry run almost covers the full velocity field. Since I fit only one geometrical parameter, I lower the number of generations in the genetic algorithm to $n_{gen} = 3$.

The distribution of the difference between recovered geometrical parameters and their input geometrical parameters are presented in figure 1.15 for the high-SNR toy models, and in figure 1.16 for the low-SNR models. Overall in the high-SNR toy models, the PA and kinematical center recovery precision is surprisingly good, given the spatial and spectral resolutions : a 1σ precision lower than ~ 3 degrees for the PA , a 1σ precision lower than 0.1 arcsec for the kinematical center for any input inclination. The quality of the adjustment of the inclination is lower, but it is a second order parameter. In intermediate inclinations regime ($40^\circ < i < 60^\circ$), the 1σ precision is close to 10° . In the case of a disk inclination of 50 degrees, a precision on the deprojected rotation curve of $\sim 30\%$ can be expected. For high input inclinations ($i = 80^\circ$), the dispersion of the distribution of the recovered inclination is large. Highest inclinations leads to a lower number of pixels available for the kinemetry expansion, and thus a lower precision. In contrast, lower inclination are also more difficult to recover, even if the number of pixels available for the kinemetry expansion is maximum. The shift of the $H\alpha$ line due to the projected rotation is lower, and thus the Gaussian fitting procedure is less precise because it is closer to the spectral resolution. It results in higher values for the non-rotation kinemetry coefficients, and makes the adjustment of the rotation term more challenging.

In the low-SNR toy models, the precision is lower as expected. No statistical bias is observed for the median difference between the input and recovered geometrical parameters. For the geometrical center, the 1σ precision is lower by a factor ~ 1.5 compared to the high-SNR toy models, at approximately 0.1 arcseconds. The 1σ precision on the PA recovery is lower by a factor ~ 2 compared to the high-SNR toy models.

Overall, this analysis demonstrates the ability to recover the geometrical parameters of an exponential disk up to a redshift 1.7 with a mean SNR lower than 10, in the area where the SNR is larger than 3. Thus, it should be possible to draw conclusions on the dynamical support of the observed galaxies in large statistical samples.

1.2.6 Towards a new kinematical classification

I have performed the fit of the geometrical parameters of MASSIV velocity fields using the KINEGEN routine and the two-step scheme described in the previous section. I have corrected the PSF effect on kinemetry ellipses flattening using the correction found in 1.2.3, with the disk scalelength evaluated from CFHT I-band morphology and the seeing measured on a star in the SINFONI λ -cube (Epinat et al., 2012). The velocity field maps of the full MASSIV sample are presented in figure 1.17. I displayed on each of them a cross and an ellipse with a semi-major axis of 1 arcsecond to illustrate the geometrical parameters recovered by the KINEGEN routine. The velocity dispersion maps are also presented in figure 1.18. Additionally, all the kinemetry expansion maps are presented in the appendix A.1.

Once the geometrical parameters are estimated, the kinematical information contained in the MASSIV sample can be interpreted using the coefficients k_n (see equation 1.5 and figures 1.19 and 1.8). To estimate the asymmetries in the kinematical data of the SINS sample, Shapiro et al. (2008) defined the coefficients v_{asym} and σ_{asym} :

$$v_{asym} = \left\langle \frac{k_{avg,v}}{B_{1,v}} \right\rangle_r, \quad (1.16)$$

$$\sigma_{asym} = \left\langle \frac{k_{avg,\sigma}}{B_{1,v}} \right\rangle_r, \quad (1.17)$$

where $k_{avg,v} = (k_{2,v} + k_{3,v} + k_{4,v} + k_{5,v})/4$ and $k_{avg,\sigma} = (k_{1,\sigma} + k_{2,\sigma} + k_{3,\sigma} + k_{4,\sigma} + k_{5,\sigma})/5$ are the sum of the asymmetric coefficients measured on the velocity and velocity dispersion fields respectively, both normalized to the rotation term $B_{1,v}$ measured on the velocity field and radially averaged to obtain v_{asym} and σ_{asym} . The indices v and σ indicate on which map (respectively radial velocity and velocity dispersion) the kinemetry coefficients were extracted. The harmonic decomposition on a velocity dispersion field is illustrated in figure 1.19. It allows to extract radially the value of the 0-th order coefficient $A_{0,\sigma}$ from which one can estimate the global velocity dispersion of the galaxy observed. In table 1.1, I display the different values measured on each velocity and velocity dispersion fields using the kinemetry expansion method.

One should note that the $k_{avg,\sigma}$ term is normalized by the rotation curve measured in the velocity field, which should, according to Shapiro et al. (2008), provide a "mass normalization" of the velocity dispersion that compensates the loss of dynamic range in external regions of the disks where the sensitivity is decreasing. Plotting both of these coefficients on a 2D diagram, Shapiro et al. (2008) set up a classification method where $K_{asym} = \sqrt{v_{asym}^2 + \sigma_{asym}^2} = 0.5$ delimitates the frontier between merger candidates ($K_{asym} > 0.5$) and isolated disk candidates

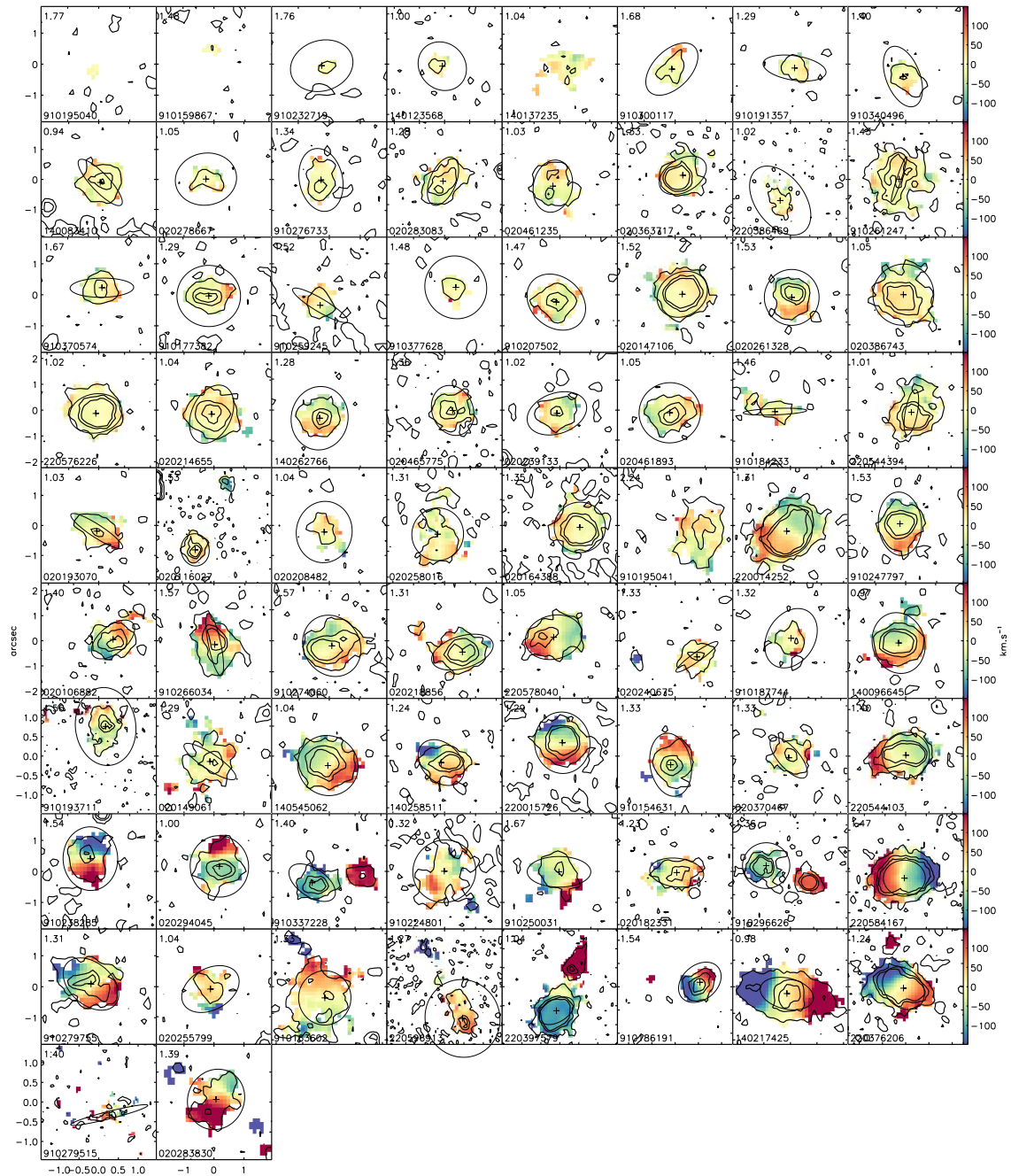


Figure 1.17 – Velocity fields of the galaxies from the MASSIV sample sorted by increasing velocity shear from top to bottom and from left to right. Contours on each maps trace the monochromatic H α flux distribution. The WDS identifier is displayed on the bottom left-hand corner of each image. Only the pixels with a signal-to-noise ratio greater than 2 are displayed. On each map we over-plotted a cross and an ellipse illustrating the recovered geometrical parameters.

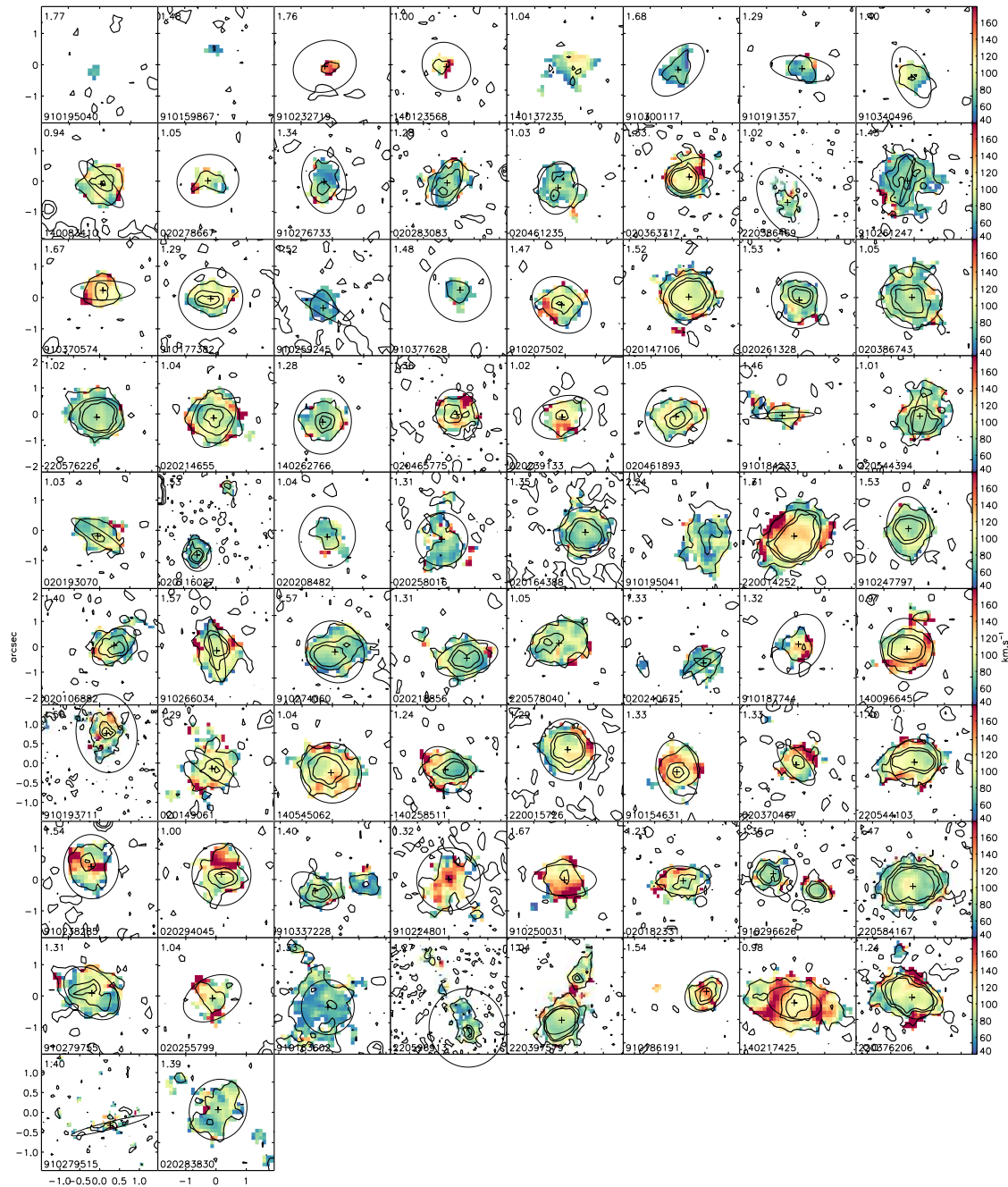


Figure 1.18 – Velocity dispersion maps of the galaxies from the MASSIV sample sorted by increasing velocity shear from top to bottom and from left to right. Contours on each maps trace the monochromatic $H\alpha$ flux distribution. The VVDS identifier is displayed on the bottom left-hand corner of each image. Only the pixels with a signal-to-noise ratio greater than 2 are displayed. On each map we over-plotted a cross and an ellipse illustrating the recovered geometrical parameters.

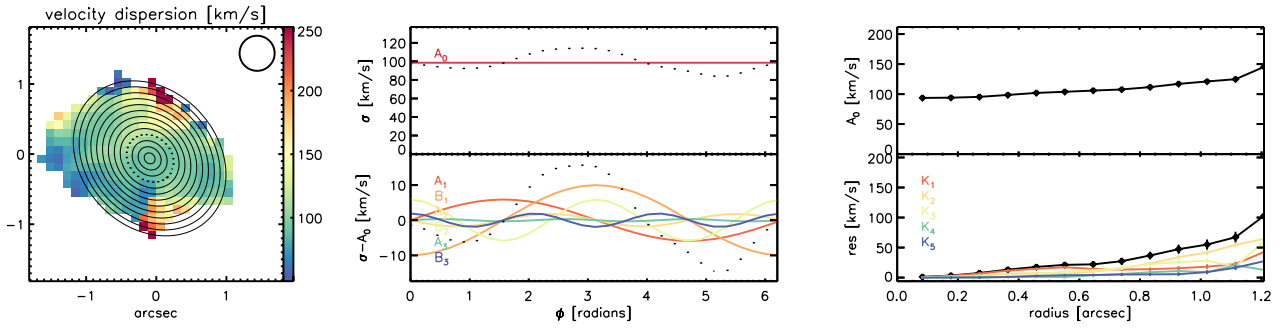


Figure 1.19 – Illustration of a kinometry run on the velocity dispersion map of the VVDS220376206 galaxy from the MASSIV sample. Left : Velocity dispersion map of the toy disk model, with kinometry ellipses over-plotted. The circle on the top right corner represents the size of the PSF (FWHM). Middle : Kinometry harmonic expansion as a function of the azimuthal angle along the ellipse with the black dots on the velocity dispersion map. The top panel shows the velocity dispersion measurements (black dots), and the fit to the kinometry velocity dispersion shift term (i.e. A_0 , red line), while the lower panel shows the velocity residuals (black dots) and the fit to the higher order harmonic coefficients. Right : Kinometry expansion for all the ellipses. The top panel shows the radial dependence of the velocity dispersion shift term (black solid line). The lower panel shows the radial dependence of the residual coefficients A_1 and K_n , $n = 2$ to 5 (see 1.5). The black line shows the quadratic sum of all the residual coefficients with an order $1 \leq n \leq 5$ (displayed using colored lines).

($K_{asym} < 0.5$). This limit was computed using a set of local observations, mergers in cosmological simulations and isolated disk toy models re-sampled to the SINFONI sampling. The original Shapiro classification diagram is presented in figure 1.20.

For each galaxy of the MASSIV sample, I measured the v_{asym} coefficient as well as the σ_{asym} coefficient for the geometrical parameters recovered with KINEGEN, that allowed me to reproduce the Shapiro et al. (2008) diagram using the MASSIV sample data. These coefficients are also estimated for the high and low SNR exponential disk toy models. The resulting diagram is presented in figure 1.21 that shows the distribution of the MASSIV sample and the disk toy models on the asymmetries diagram.

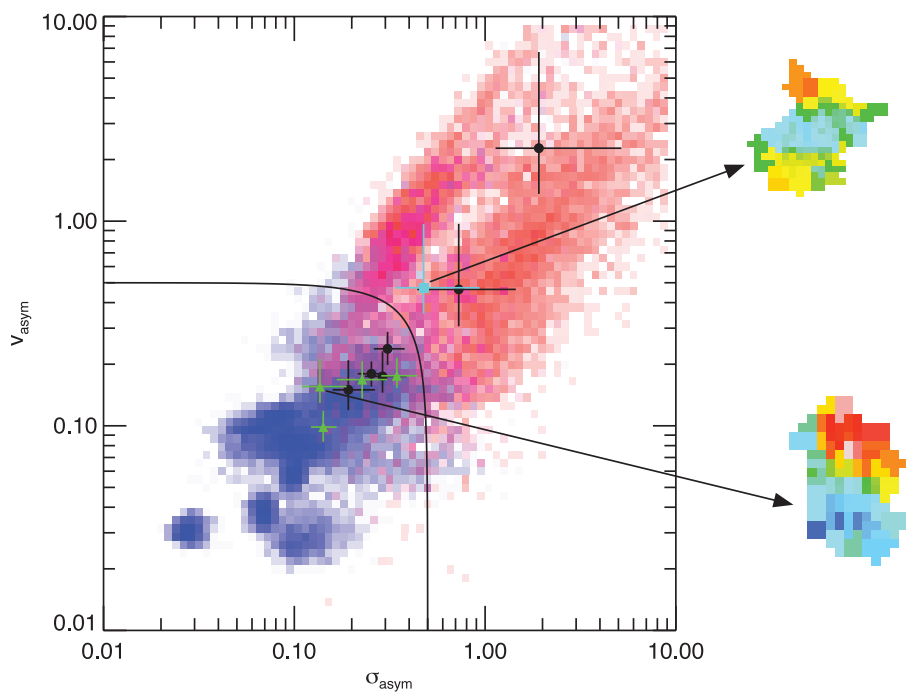


Figure 1.20 – Asymmetry measures using kinemetry performed on the SINS sample. The blue pixels trace the distribution of the isolated disks simulated velocity fields, while the red pixels trace the distribution of the merging galaxies simulated velocity fields. The black line shows the limit $K_{asym} = 0.5$. Galaxies above this line are classified as mergers, while galaxies below this line are classified as isolated disk. The position of the SINS sample is over-plotted, and an example for the isolated disk and merger classification is showed. Credits : Shapiro et al. (2008)

Table 1.1 – Full MASSIV sample kinemetry parameters recovered from the kinemetry expansion. From left to right : 1. MASSIV VVDS identifier. 2. Redshift. 3. PSF full width at half maximum. 4. Inclination recovered from KINEGEN. 5. Maximum velocity measured in the $B_{1,v}$ coefficient. 6. Mean SNR with line detection above 3σ . 7. Ratio optical radius over seeing FWHM. 8. Radial velocity asymmetry coefficient. 9. Velocity dispersion asymmetry coefficient. 10. Mean value for the velocity dispersion measured in the $A_{0,\sigma}$.

ID	z	seeing [arcsec]	i [degree]	v_{max} [km.s ⁻¹]	SNR	$r_{last}/seeing$	v_{asym}	σ_{asym}	σ [km.s ⁻¹]
020106882	1.40	0.49	57.82	222.46	7.60	1.20	0.13	0.18	104.94
020116027	1.53	0.60	29.81	25.60	8.20	1.30	0.23	0.31	67.67
020147106	1.52	0.65	34.94	14.20	12.20	1.40	0.35	0.52	99.02
020149061	1.29	0.85	75.01	51.34	6.20	0.70	0.36	0.67	92.29
020164388	1.35	0.83	10.21	27.70	11.00	1.20	0.23	0.29	78.28
020182331	1.23	0.74	58.06	66.71	6.10	0.90	0.19	0.26	102.51
020193070	1.03	0.58	77.97	60.61	4.90	0.80	0.12	0.19	82.38
020208482	1.04	0.58	24.96	26.84	4.10	0.30	0.06	0.15	75.99
020214655	1.04	0.87	49.91	30.88	7.50	0.80	0.27	0.49	93.82
020218856	1.31	0.56	59.68	46.95	8.10	1.30	0.15	0.19	81.74
020239133	1.02	0.79	48.87	51.63	5.30	0.80	0.19	0.55	117.11
020240675	1.33	0.85	79.78	11.96	5.50	0.50	1.75	0.75	68.87
020255799	1.04	0.76	47.77	25.41	4.50	0.60	0.39	0.58	100.53
020258016	1.31	0.60	11.76	27.38	4.70	1.40	0.93	0.73	74.96
020261328	1.53	0.62	11.71	89.03	7.80	1.00	0.13	0.18	90.63
020278667	1.05	0.65	32.28	10.85	4.10	0.20	0.63	1.10	102.37
020283083	1.28	0.78	59.39	31.70	7.00	0.90	0.19	0.31	74.85
020283830	1.39	0.77	24.12	118.47	4.80	1.20	0.07	0.09	81.67
020294045	1.00	0.59	18.41	99.14	7.00	1.20	0.23	0.40	117.44
020363717	1.33	0.64	50.20	22.07	11.40	1.10	0.79	1.31	112.75
020370467	1.33	0.71	65.91	27.55	5.50	0.90	0.77	1.23	102.40
020386743	1.05	0.73	10.64	31.83	8.80	0.90	0.31	0.54	106.53
020461235	1.03	0.63	42.51	42.13	5.80	1.10	0.09	0.13	76.05
020461893	1.05	0.60	11.91	39.34	7.00	1.30	0.12	0.28	94.93
020465775	1.36	0.88	36.75	28.28	6.80	0.70	0.59	0.92	110.75
140083410	0.94	0.69	73.13	18.76	6.00	1.00	0.41	0.47	95.13
140096645	0.97	0.56	31.83	75.88	10.30	1.00	0.06	0.14	115.24
140137235	1.04	0.76	62.00	33.98	3.20	0.10	0.17	0.33	–
140217425	0.98	0.95	55.29	289.27	6.80	1.90	0.09	0.14	142.40
140258511	1.24	0.49	43.95	199.71	8.10	1.30	0.14	0.17	102.43
140262766	1.28	0.51	26.72	57.13	6.50	1.00	0.08	0.09	80.48
140545062	1.04	0.70	16.25	89.63	8.50	1.30	0.06	0.15	108.40
220014252	1.31	0.70	25.48	83.05	11.00	1.70	0.30	0.22	122.76
220015726	1.29	0.46	25.81	158.14	10.40	1.00	0.09	0.08	104.51
220376206	1.24	0.50	39.07	148.26	12.60	2.40	0.08	0.12	109.87
220386469	1.02	0.23	51.25	10.39	4.50	1.10	1.48	3.22	81.12
220397579	1.04	0.64	53.82	34.33	15.70	2.00	0.54	0.66	87.94
220544103	1.40	0.76	14.02	58.73	10.40	1.20	0.13	0.23	101.42
220544394	1.01	0.58	74.64	30.83	10.50	1.10	0.21	1.45	76.83
220576226	1.02	0.58	42.24	18.96	11.60	1.30	0.20	0.26	79.48
220578040	1.05	0.62	40.81	102.22	5.90	1.40	0.11	0.16	81.43
220584167	1.47	0.75	31.00	171.00	13.80	2.10	0.06	0.07	90.05
220596913	1.27	0.18	17.55	27.63	5.70	6.20	0.28	0.39	72.54

End of table 1.1. Full MASSIV sample kinemetry parameters recovered from the kinemetry expansion.

ID	z	Seeing [arcsec]	i [degree]	v_{max} [km.s ⁻¹]	SNR	$r_{last}/seeing$	v_{asym}	σ_{asym}	σ [km.s ⁻¹]
910154631	1.33	0.56	49.90	69.85	6.70	1.10	0.20	0.22	118.82
910163602	1.33	0.59	15.33	192.89	5.00	2.50	0.16	0.15	75.32
910177382	1.29	0.52	15.87	21.77	7.30	1.30	0.19	0.27	101.18
910184233	1.46	0.61	84.21	9.51	3.90	0.30	0.90	1.14	84.23
910186191	1.54	0.47	42.90	214.60	–	–	0.06	0.11	134.73
910187744	1.32	0.75	42.03	19.59	5.10	0.40	0.26	0.79	111.96
910191357	1.29	0.58	67.42	15.21	3.80	0.40	0.17	0.45	69.23
910193711	1.56	0.27	39.43	45.15	6.00	1.80	0.33	0.73	99.80
910207502	1.47	0.63	37.69	19.48	6.20	0.80	0.46	0.59	100.28
910224801	1.32	0.49	13.54	64.13	4.40	1.80	0.37	0.41	133.98
910232719	1.76	0.39	37.76	64.13	–	–	0.37	0.41	133.98
910238285	1.54	0.47	30.87	158.24	5.70	1.30	0.05	0.11	120.65
910247797	1.53	0.30	47.25	81.92	9.40	2.70	0.25	0.26	77.31
910250031	1.67	0.81	61.24	41.77	5.10	0.70	0.29	0.55	130.68
910259245	1.52	0.73	78.06	44.13	4.30	0.70	0.05	0.08	57.81
910261247	1.43	0.62	84.47	78.09	8.20	1.70	0.69	0.51	70.89
910266034	1.57	0.55	78.76	135.21	7.50	1.60	0.06	0.09	92.91
910274060	1.57	0.87	28.70	61.55	9.50	0.80	0.10	0.12	79.72
910276733	1.34	0.37	47.22	15.29	5.70	0.90	0.23	0.52	66.49
910279515	1.40	0.21	84.64	126.51	3.50	1.50	0.10	0.08	82.09
910279755	1.31	0.58	48.46	314.64	6.90	1.60	0.12	0.19	86.27
910296626	1.36	0.53	14.93	314.64	10.90	0.60	0.12	0.19	86.27
910300117	1.68	0.68	53.98	20.49	4.10	0.80	0.26	0.32	67.32
910337228	1.40	0.76	57.65	21.37	7.50	0.70	0.98	1.73	94.97
910340496	1.40	0.47	57.85	35.91	5.40	0.80	0.07	0.22	90.32
910370574	1.67	0.55	72.43	25.20	6.10	0.80	0.12	0.17	120.13
910377628	1.48	0.69	16.16	6.59	4.10	0.30	0.28	0.67	76.47

In the MASSIV sample kinematical data, we find 43 galaxies below $K_{asym} = 0.5$ and eligible for a “isolated disk” classification, and 28 galaxies above this limit, eligible for a “galaxy merger” classification. Almost 40% of the galaxies in the MASSIV sample where the H α (or [OII]) is detected and spatially resolved fall in the merger region. The Shapiro classification on the v_{asym}/σ_{asym} diagram provides a very similar information to the classification of Epinat et al. (2012) (see figure 1.22). Indeed, in figure 1.21, I plotted the MASSIV galaxies using symbols that indicate the classification scheme of the first epoch MASSIV sample, extended in my work to the full MASSIV sample. Almost all the objects classified as rotating disks (28/29) are classified as isolated disks on the Shapiro diagram, and objects classified as non-rotating disks have a high probability to be classified as mergers (22/37) on the Shapiro diagram. In my analysis, I find that the residuals to the model of Epinat et al. (2012) are strongly correlated to the v_{asym} coefficient, confirming the correlation between the two classifications.

The dispersion of the v_{asym} and σ_{asym} coefficients is somewhat surprising because the low SNR and high SNR disk toy models are built with the exact same rotation model : one could expect all the coefficients v_{asym} and σ_{asym} to fall on a very restricted area in the Shapiro diagram. There is two explanations for this dispersion :

- The toy models are projected at different redshifts, and they are inserted in a real SINFONI λ cube at different positions in the spectra. Some redshifts are more affected by OH sky lines, which increases the perceived amount of disturbance in the velocity field.

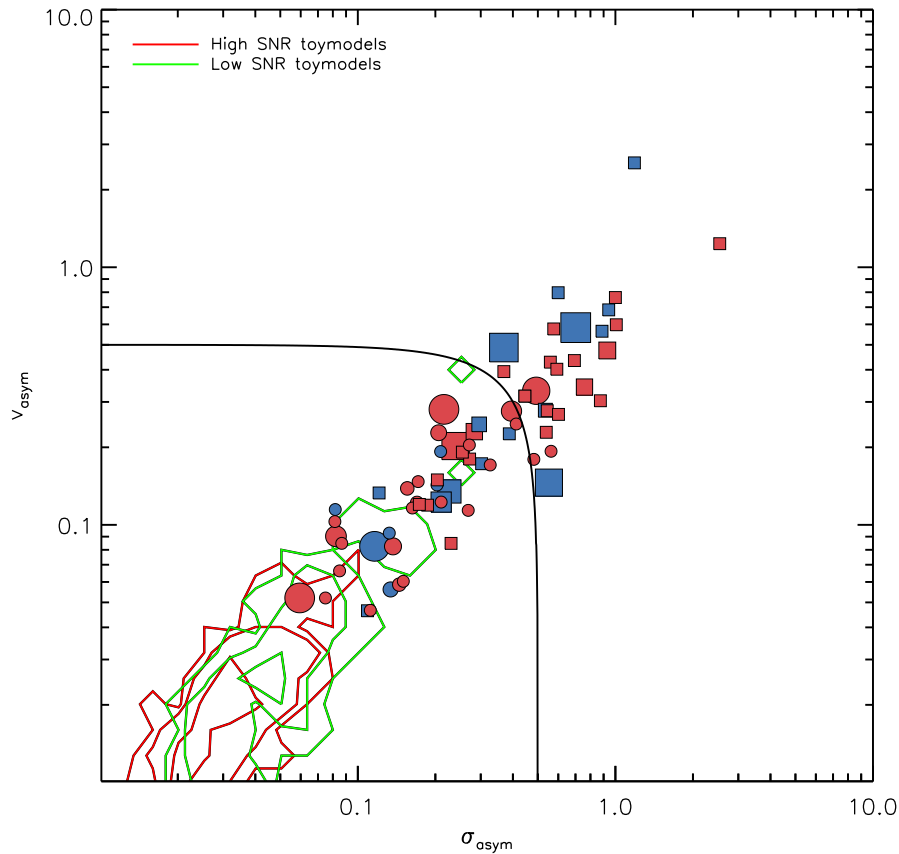


Figure 1.21 – Asymmetry measures using kinemetry performed on the MASSIV sample and the disk toy models presented in section 1.2.5. Squares and circles symbols represent the position of MASSIV galaxies on the diagram. The squares and circles indicate respectively the classification of “rotating” and “non-rotating” objects obtained in Epinat et al. (2012). Blue and red symbols represent respectively galaxies classified as interacting and isolated. The size of the symbols is related to the SNR of the observation. The black line shows the limit $K_{asym} = 0.5$. There are 43 galaxies below this limit (candidates for isolated disk classification) and 28 galaxies above this limit (candidates for merger classification), i.e. $\sim 40\%$ of galaxies eligible for merger classification. The two colored contours are displaying the position of the high and low SNR exponential disk toy models on the diagram.

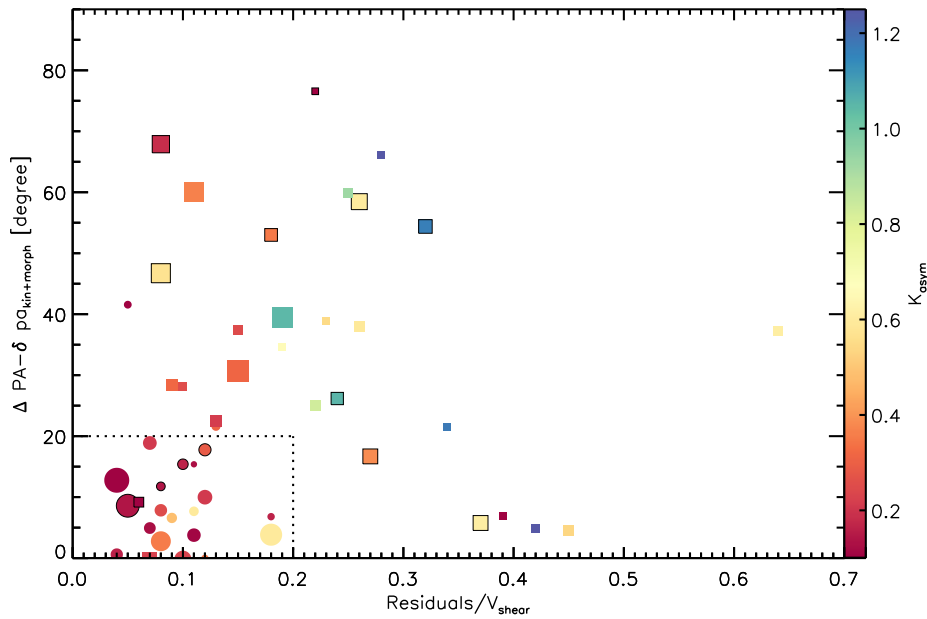


Figure 1.22 – Disagreement between morphological and kinematic position angles as a function of the velocity field residuals normalized by the velocity shear, as in Epinat et al. (2012) but using the full MASSIV sample. Dots and squares represent rotating and non-rotating galaxies. Symbols with black contours mark interacting galaxies. The color of the symbols encodes the value of the kinemetry coefficient K_{asym} . The symbol size is related to the mean SNR of the SINFONI data. The lower-left region drawn by dotted lines delimitates the “rotating disk” region.

- The construction of the coefficients v_{asym} and σ_{asym} using a normalization to the rotation curve naturally introduces dispersion when the inclination of the toy model varies because the rotation is projected on the plane of the sky.

Furthermore, the strong correlation between v_{asym} with σ_{asym} implies that the normalization with the rotation term $B_{1,v}$ dominates the dynamical range of the coefficients. Such a strong correlation demonstrates that one coefficient is sufficient to provide the exact same classification. The classification frontier $K_{asym} = 0.5$ is calibrated on very different models (toy models, projected observations, and cosmological simulations) that are to my interpretation very difficult to compare : (i) the mergers cosmological simulations lack the resolution to resolve numerically the vertical structure of gas disk, (ii) the local observed galaxies have gas fractions much lower than young galaxies and (iii) rotating disk toy models does not provide any modeling of the motions other than pure rotation. It is clear from the kinemetry analysis of the exponential disk toy models presented in section 1.2.5 that high redshift galaxies are clearly not only dynamically supported by rotation. Both high and low SNR toy models samples fall in the low left corner of the Shapiro kinemetry diagram, meaning that the harmonic decomposition is dominated by the B_1 term. Only the toy models which falls close to OH sky line in the SINFONI spectra display asymmetries comparable to the best case of rotating disk in the MASSIV sample. However, tracing frontiers between two populations of mergers and isolated disks is obviously not reflecting the complexity of the processes leading to strongly disturbed kinematics. Galaxies dynamically perturbed are relaxing their gas

component over hundreds of million years, and the disturbance of the gaseous component slowly decreases until it reaches the most stable state for the system.

At this point, I considered that it remains difficult to determine unambiguous estimators that allow to distinguish recent or ongoing mergers from isolated rotating disks in spectroscopic data with spatial and spectral resolution comparable to the one of the SINFONI instrument. The kinematical signatures of a high redshift merger with a high gas fraction is unclear because isolated disks in this redshift range are expected to experience a strong turbulence and gravitational fragmentation (Bournaud et Elmegreen, 2009) due to the elevated gas fractions Daddi et al. (2010a). These violent instabilities may make the distinction between isolated disks from mergers difficult. In order to address the ability to detect recent or on-going merger kinematical signatures using the kinemetry analysis, I need an homogeneous sample of numerical simulations of high redshift mergers and isolated disks. This sample should provide a numerical resolution able to capture the highly complex behavior of the gaseous disk. Finally, I need this simulation sample to statistically probe the most important parameters in terms of impact on the gas kinematics, to constrain mergers on a specific area of the disturbance estimators space. In the following chapters, I will show how I built such sample of simulations, and what informations I can extract in terms of kinematical classification.

Conditions Initiales Idéalisées

2

Résumé

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Dans le précédent chapitre, nous avons vu la nécessité de créer un échantillon de simulations numériques afin de caractériser les signatures de fusions à grand décalage spectral. Créer un tel échantillon de simulations nécessite de contrôler les paramètres qui le définissent, i.e. les paramètres physiques de modèles numériques de galaxies. Ce chapitre traite donc de la création de modèles idéalisés de galaxies en équilibre dynamique qui nous permettront de faire avancer notre analyse cinématique de l'échantillon MASSIV. La méthode numérique présentée ne se restreint pas au seul cas des modèles de galaxies à grand décalage spectral, et peut être appliquée à n'importe quel cas de figure de par sa versatilité.

Dans le cadre de cette thèse, j'ai souhaité réaliser plusieurs simulations de fusions de disques dont je puisse contrôler tous les paramètres (masse, taille, orbite, etc.) afin de pouvoir contraindre au maximum le comportement du gaz dans le contexte des galaxies à grand décalage spectral. Je me suis donc tourné vers les simulations dites "idéalisées", c'est-à-dire des simulations qui s'effectuent dans une boîte fermée sans échange possible avec le milieu extérieur pendant la durée de la simulation. Le contexte cosmologique sert cependant à déterminer les conditions initiales au lancement de la simulation. Les simulations idéalisées permettent une paramétrisation plus complète des systèmes de galaxies que l'on souhaite étudier que les simulations cosmologiques. Mais encore faut-il disposer de modèles de galaxies dont la structure reste à l'équilibre lors de l'intégration temporelle... Dans les sections suivantes, je décris une méthode de construction de conditions initiales dynamiquement stables, qui suit globalement la méthode proposée dans Springel et al. (2005b).

2.1 Positionnement des particules

La première étape dans la construction d'un modèle idéalisé de galaxie dynamiquement stable passe par la réalisation d'un positionnement de particules Lagrangiennes qui permettent de décrire un profil de densité volumique prédéfini. Si la fonction de répartition d'une loi de probabilité est connue analytiquement, il est alors facile de réaliser un tirage aléatoire qui suivra cette loi. Néanmoins, le calcul de la fonction de répartition implique des intégrations qui ne sont parfois pas résolubles analytiquement, ce qui restreint fortement le choix de la loi de distribution à des cas où la fonction de répartition est connue analytiquement. Plusieurs algorithmes permettent de construire des réalisations statistiques sans passer par la fonction de répartition, ce qui permet de ne pas être contraint dans le choix des profils de densité initiaux du modèle numérique. L'algorithme Monte Carlo Markov Chain (MCMC) permet de se passer de la fonction de répartition pour construire une distribution suivant une loi de probabilité (i.e. une loi de densité). Mon choix s'est porté sur l'implémentation de Metropolis et al. (1953) qui implique une marche aléatoire dans l'espace des paramètres de la fonction de densité (espace cartésien 3D dans le cas présent) dont la probabilité de déplacement d'un point à l'autre de l'espace des paramètres est Gaussienne. La construction de la chaîne de Markov se déroule ainsi :

1. Initialisation de la chaîne, positionnement de la première particule Lagrangienne à $x_{i=0}$, avec x_i la coordonnée Cartésienne de la particule i de la chaîne MCMC.
2. Génération d'une position candidate x_p pour la particule suivant x_i , avec $x_p = x_i + \sigma$ et σ tiré aléatoirement suivant une loi Gaussienne.
3. Estimation de la probabilité $P(x_p, x_i)$ de la position candidate x_p , avec $P(x_p, x_i) = \frac{\rho(x_i)p(x_i, x_p)}{\rho(x_p)p(x_p, x_i)}$ si $\rho(x_i)p(x_i, x_p) < \rho(x_p)p(x_p, x_i)$ et $P(x_p, x_i) = 1$ sinon. $\rho(x_i)$ est la densité aux coordonnées x_i de la particule i , et $p(x_i, x_p)$ est la probabilité de déplacement du point x_i au point x_p . Dans le cas de l'implémentation de Metropolis-Hasting la loi de déplacement aléatoire est une Gaussienne de largeur fixe quelle que soit la position considérée, donc $p(x_i, x_p) = p(x_p, x_i)$. Dans des formulations alternatives à Metropolis et al. (1953), la loi de déplacement peut être fonction de la position dans l'espace des paramètres de la loi de probabilité reconstruite par la chaîne.
4. Tirage aléatoire uniforme dans $[0, 1]$ stocké dans α
5. Si $\alpha \leq P(x_p, x_i)$ alors $x_{i+1} = x_p$, sinon $x_{i+1} = x_i$ i.e. la position de la particule est identique à la précédente.
6. Répéter à partir de l'étape (2) jusqu'à ce que la distribution atteigne le nombre de particules voulu.

Cet algorithme nécessite l'utilisation intensive d'un générateur de nombres aléatoires, afin d'assurer une convergence rapide vers la fonction de densité cible. L'intérêt de l'algorithme MCMC réside dans l'estimation de la probabilité d'existence d'une position candidate, qui repose sur un rapport de la fonction de densité volumique (point 3 de la liste ci-dessus). Nul besoin donc d'avoir à disposition la fonction de répartition, ce qui simplifie grandement la réalisation de la distribution de particules Lagrangiennes. Divers test numériques suggèrent de définir une dispersion σ_{jump} proportionnelle à l'échelle de longueur de la fonction de densité de la composante considérée pour la loi de proposition de déplacement Gaussienne. La qualité de la convergence de l'algorithme peut être estimée par le taux d'acceptance $\mathcal{A} = n_{accept}/n_{chain}$ dans la chaîne Monte-Carlo, qui est le rapport du nombre de

propositions de déplacement acceptées n_{accept} sur le nombre d'éléments de la chaîne n_{chain} . Une acceptante faible signifie que la loi de déplacement Gaussienne propose des positions systématiquement éloignées du centre de masse de la composante, tandis qu'un taux d'acceptante élevé signifie que la chaîne n'atteint pas les régions éloignées du centre de masse de la composante. Numériquement, j'ai estimé l'acceptance idéale à 75%. Pour $\mathcal{A} > 90\%$, il est nécessaire d'augmenter σ_{jump} . Inversement, pour $\mathcal{A} < 50\%$, il est nécessaire de diminuer σ_{jump} . Pour les structures en disque de manière générale, la dispersion idéale de la loi Gaussienne de déplacement est égale à la moitié de la longueur d'échelle de la structure. Pour les halos qui sont des structures plus étendues de manière générale, la dispersion idéale de la loi Gaussienne de déplacement est égale au quart de la longueur d'échelle de la structure. Pour chaque composante de la galaxie (disque stellaire, disque gazeux, bulbe stellaire, halo de matière sombre), nous pouvons donc spécifier une fonction de densité volumique qui permettra de réaliser une distribution aléatoire de particules Lagrangiennes suivant ce profil. Pour le disque stellaire et gazeux, le profil de densité le plus communément utilisé est le profil exponentiel, exprimé dans un référentiel cylindrique :

$$\rho(r, z) = \frac{M}{2\pi z_s^2} \exp(-r/r_s) \operatorname{sech}(-z/z_s), \quad (2.1)$$

avec M la masse totale de la composante, r_s l'échelle de longueur radiale du disque, z_s l'échelle de longueur azimutale du disque. Le profil de densité de Miyamoto et Nagai (1975), analytiquement plus complexe, est lui aussi fréquemment utilisé. Il est décrit par l'expression

$$\rho(r, z) = \left(\frac{M}{4\pi} \right) \frac{r_s r^2 + (r_s + 3\sqrt{z^2 + z_s^2})(r_s + \sqrt{z^2 + z_s^2})}{[r^2 + (r_s + \sqrt{z^2 + z_s^2})^2]^{5/2} (z^2 + z_s^2)^{3/2}}. \quad (2.2)$$

Le halo de matière sombre ainsi que le bulbe stellaire sont quant à eux généralement modélisés par un système sphérique ou ovoïdal. Les simulations cosmologiques N-corps forment des halos de matière sombre avec un profil analytiquement décrit par Navarro et al. (1997) :

$$\rho(r) = \frac{\rho_0}{\frac{r}{r_s} \left(1 + \frac{r}{r_s}\right)^2}. \quad (2.3)$$

avec r_s l'échelle de longueur du halo, et ρ_0 la densité centrale du halo. Le paramètre de concentration du halo NFW est défini par la fraction :

$$c = \frac{r_{200}}{r_s} \quad (2.4)$$

avec r_{200} le rayon de Viriel. Dans les simulations N-corps cosmologiques, le paramètre de concentration du halo de matière sombre pour les galaxies de type Voie Lactée ($v_{200} \sim 160$ km.s⁻¹), c est compris entre 10 et 15. Les halos de matière noire sont aussi fréquemment modélisés avec un profil de Hernquist (1990)

$$\rho(r) = \frac{M}{2\pi} \frac{a}{r(r+a)^3}, \quad (2.5)$$

avec a une longueur d'échelle que l'on peut relier au paramètre de concentration du profil de NFW en imposant une densité centrale égale entre les deux profils :

$$a = r_{halo} \sqrt{2 \left(\ln(1+c) - \frac{c}{1+c} \right)}. \quad (2.6)$$

Le profil de Hernquist possède une pente dans les régions centrales ($r < 0.1r/r_{200}$) similaire à celle du profil NFW, mais est caractérisé par une décroissance radiale plus prononcée au delà de ce rayon. Ce profil, que j'ai utilisé fréquemment dans mon travail sur les conditions initiales, est plus facile à manipuler dans le sens où il fournit une coupure en densité plus prononcée que le profil NFW ; permettant de construire un halo dans un volume fini. Le nombre de particules par unité de volume déterminant la résolution numérique, la taille du volume détermine par extension la résolution numérique du modèle de galaxie considéré.

La convergence de l'algorithme MCMC Metropolis Hasting est illustrée dans la figure 2.1 pour le profil de densité de Miyamoto et Nagai (1975) en fonction du nombre de particules Lagrangiennes utilisées. La figure 2.2 permet de visualiser "l'empilement" des différentes composantes pour un modèle numérique de galaxie à disque.

2.2 Potentiel gravitationnel

L'étape suivant la création d'une distribution de particules Lagrangiennes est le calcul du potentiel gravitationnel de la distribution. Pour ce faire, il est nécessaire d'interpoler la valeur de la densité volumique sur une grille cartésienne 3D dont le nombre d'éléments définit la précision numérique sur le calcul de cette valeur. Le potentiel gravitationnel $\Phi(\mathbf{x})$ est alors calculé en intégrant l'équation de Poisson :

$$\Phi(\mathbf{x}) = \int \mathcal{G}(\mathbf{x}, \mathbf{x}') 4\pi\rho(\mathbf{x}') d^3\mathbf{x}', \quad (2.7)$$

avec x le vecteur position et \mathcal{G} la fonction de Green. Cette équation peut s'interpréter comme une convolution 3D de la fonction de densité volumique par la fonction de Green. Cette opération est donc effectuée en passant dans l'espace de Fourier où une convolution correspond au produit des transformées de Fourier des deux fonctions :

$$\Phi(\mathbf{x}) = \mathcal{F}^{-1}(\mathcal{F}(4\pi\rho(\mathbf{x})) \otimes \mathcal{F}(\mathcal{G}(\mathbf{x}))) \quad (2.8)$$

où \mathcal{F} et \mathcal{F}^{-1} sont respectivement les transformées de Fourier directe et inverse.

Cette méthode a pour avantage d'être très rapide grâce aux implémentations performantes de l'algorithme Fast Fourier Transform (FFT), mais a pour inconvénient d'avoir une approche périodique du problème. Ce problème nécessite donc de placer les modèles de galaxies le plus au centre possible de la grille cartésienne de potentiel. Si le calcul est réalisé en l'état, le potentiel au bord de la grille va donc être affecté par le potentiel à l'autre extrémité de la grille, et vice-versa. Afin de pallier à cette difficulté, nous choisissons d'utiliser la méthode du "zero-padding" décrite dans Hockney et Eastwood (1988), qui consiste à étendre chaque dimension de la grille cartésienne d'un facteur deux. Cette extension permet de concentrer les effets de périodicité liés au passage dans le domaine de Fourier dans les sept quadrants supplémentaires de la grille. Bien que rapide, l'implémentation numérique de cette méthode nécessite de réserver une grande quantité de mémoire vive lors de l'implémentation numérique, puisque pour estimer le potentiel sur une grille de 256^3 éléments, il est nécessaire de réserver en mémoire une grille de 512^3 éléments.

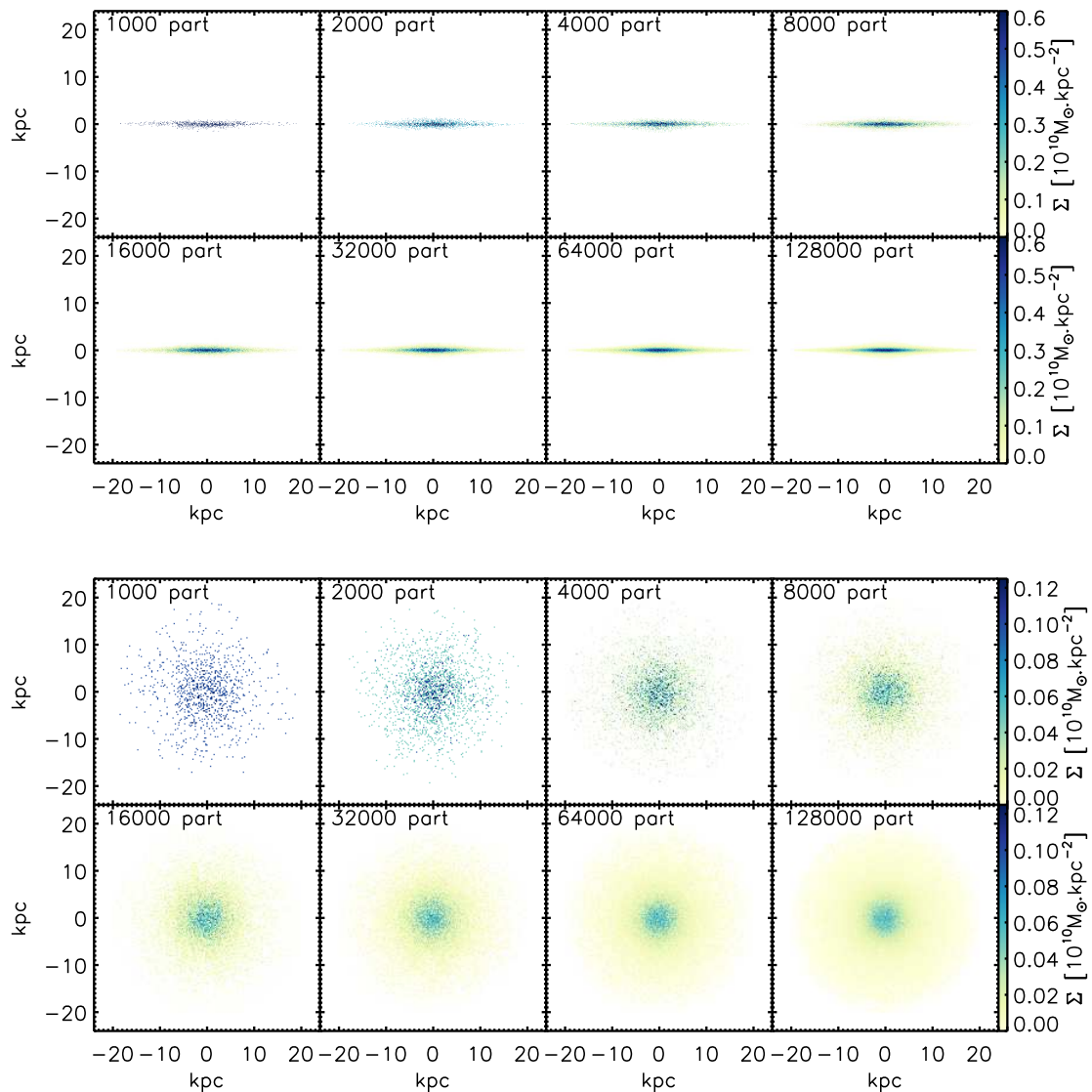


Figure 2.1 – Convergence de l’algorithme MCMC Metropolis Hasting pour un disque de masse totale $M = 6.15 \times 10^{10} M_{\odot}$ avec un profil de densité volumique de Miyamoto et Nagai (1975), une longueur d’échelle radiale $r_d=5\text{kpc}$ et une longueur d’échelle verticale $h_d=500\text{ pc}$. La distribution de Miyamoto et Nagai est coupée pour la densité volumique $\rho_{min} = \rho(r = r_{cut}, z = 0)$ avec $r_{cut}=20\text{ kpc}$. Huit disques sont construits de manière identique, mais avec un nombre total de particules Lagrangiennes variant entre 1000 et 128000. En haut : densité surfacique des disques vus par la tranche. En bas : densité surfacique des disques vus de face. Les disques construits avec moins de 8000 particules montrent une densité surfacique bruitée et éloignée du profil d’entrée. La qualité de la convergence de l’algorithme MCMC commence à permettre une reconstruction satisfaisante de la fonction de densité pour un nombre de particules supérieur ou égal à 16000.

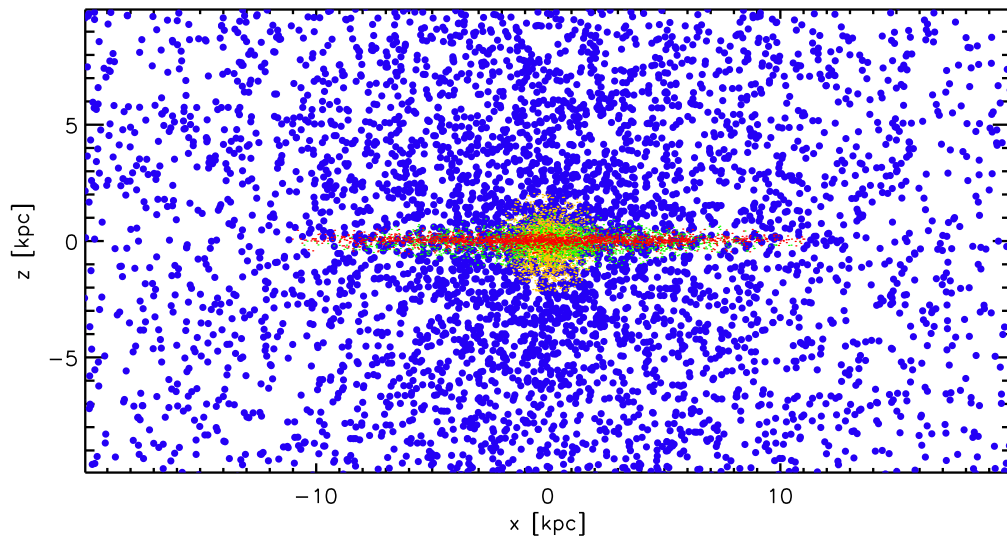


Figure 2.2 – Distribution spatiale d'un modèle idéalisé de galaxie avec un halo (points bleus), un bulbe stellaire (points oranges), un disque stellaire (points verts) et un disque de gaz (points rouges). Chacune des composantes est construite avec l'algorithme MCMC avec 10^3 particules. Le halo et le bulbe sont modélisés avec un profil de Hernquist (1990), tandis que le disque stellaire et gazeux sont modélisés avec un profil de Miyamoto et Nagai (1975). Chaque point représente une particule Lagrangienne dont la masse est proportionnelle à la taille du point sur le graphique.

2.3 Equilibre hydrostatique

La construction d'un modèle de galaxie à l'équilibre nécessite une étape spécifique pour la mise à l'équilibre hydrostatique du gaz. En effet, notre définition initiale de la fonction de densité volumique du gaz ne suppose pas cet équilibre. Afin de pouvoir obtenir cet équilibre, nous supposons que le gaz se trouve dans une structure verticalement fine. D'après l'équation d'Euler pour des fluides parfaits, l'équilibre hydrostatique est réalisé pour la condition :

$$\frac{1}{\rho_{\text{gaz}}(r, z)} \frac{\partial P(r, z)}{\partial z} + \frac{\partial \Phi(r, z)}{\partial z} = 0, \quad (2.9)$$

avec $P(r, z)$ la pression, $\Phi(r, z)$ le terme de potentiel gravitationnel et $\rho_{\text{gaz}}(r, z)$ la densité volumique du gaz. Dans le cas d'un gaz isotherme, le terme de pression s'écrit :

$$P(r, z) = c_s^2 \rho_{\text{gaz}}(r, z), \quad (2.10)$$

avec c_s la vitesse du son dans le gaz. L'intégration de l'équation 2.9 permet d'obtenir une expression de la densité du gaz en fonction du gradient de potentiel gravitationnel :

$$\rho_{\text{gaz}}(r, z) = \rho_{\text{gaz}}(r, z = 0) \exp\left(-\frac{\Delta\Phi_z(r, z)}{c_s^2}\right), \quad (2.11)$$

avec $\Delta\Phi_z(r, z) = \Phi(r, z) - \Phi(r, z = 0)$ la différence verticale de potentiel gravitationnel entre le plan équatorial et l'altitude considérée z . La densité de gaz dans le plan équatorial de la galaxie reste inconnue à ce stade de la méthode. Afin de réaliser l'équilibre hydrostatique vertical du disque, nous initialisons la densité dans le plan équatorial avec $\rho_0 = \Sigma_{\text{gas}}(r)$ avec $\Sigma_{\text{gas}}(r)$ la densité surfacique du gaz. Il devient alors possible de faire converger de manière itérative la structure verticale du disque vers un état d'équilibre hydrostatique. En ré-utilisant l'algorithme Metropolis-Hasting précédemment décrit, nous re-calculons l'altitude z des particules Lagrangiennes modélisant le gaz en suivant l'équation 2.11. Chaque modification de la structure du disque implique une ré-évaluation complète du potentiel gravitationnel. En répétant l'opération une dizaine de fois, il est possible de converger vers une structure verticale du disque de gaz. La procédure itérative de mise à l'équilibre hydrostatique est réalisée en suivant les instructions suivantes :

1. Initialisation de la structure du disque de gaz avec une fonction de densité initiale.
2. Calcul du potentiel gravitationnel.
3. Calcul de la densité du gaz dans le plan équatorial.
4. Modification de la hauteur z de chaque particule de gaz de manière à vérifier l'équation

$$\rho_{\text{gaz}}(r, z) = \rho_{\text{gaz}}(r, z = 0) \exp\left(-\frac{\Delta\Phi_z(r, z)}{c_s^2}\right)$$

5. Nouvelle estimation de la masse du système en tenant compte des critères de coupure de la fonction de densité.
6. Retour au point (2) jusqu'à la convergence de la structure verticale du disque.

Pour des températures de gaz relativement basses ($T < 10^4 \text{K}$), la structure verticale du disque peut devenir très fine. Le choix de la résolution numérique de la grille permettant d'évaluer le potentiel gravitationnel est donc critique car une résolution trop faible amènerait à sur-estimer l'épaisseur du disque et pourrait éloigner le disque de l'état d'équilibre hydrostatique recherché.

2.4 Calcul des vitesses

Afin de compléter la description de notre modèle numérique de galaxie, nous devons spécifier les vitesses Cartésiennes de chaque particule Lagrangienne. Pour qu'une particule reste sur une orbite circulaire, sa vitesse doit être définie par :

$$v_c(r) = r \frac{\partial \Phi(r)}{\partial r}, \quad (2.12)$$

avec v_c la vitesse circulaire au rayon r . La figure 2.3 montre les différentes courbes de rotation obtenues par le calcul de v_c pour chacune des composantes d'un modèle idéalisé de galaxie jeune.

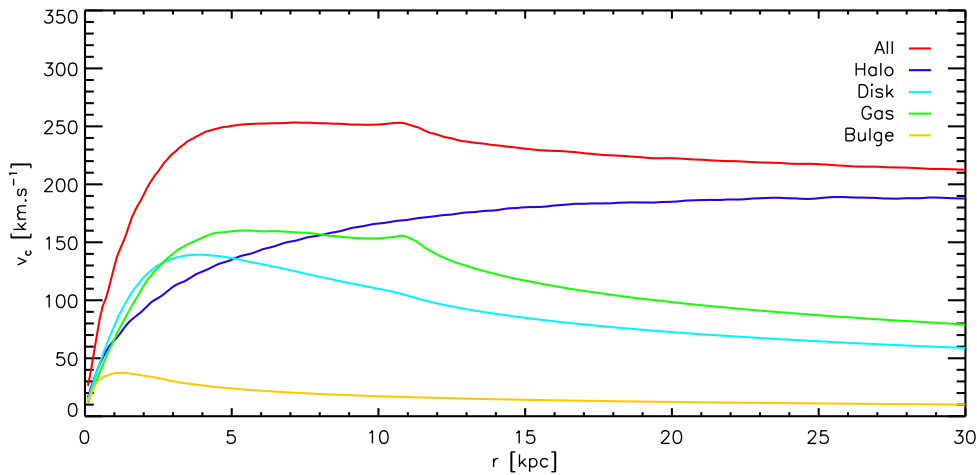


Figure 2.3 – Courbe de rotation pour un modèle de galaxie typique de l'Univers à $z=2$ avec une vitesse de Viriel $V_{200}=210 \text{ km.s}^{-1}$ calculée à partir de la dérivée numérique du potentiel gravitationnel. La masse baryonique dans le disque stellaire représente une fraction $m_d = 0.1M_{200}$ de la masse de Viriel, la fraction de masse du bulbe sur la masse du disque est $B/D = 0.04$ et le disque de gaz contient une fraction $f_g = 0.65$ de la masse totale baryonique située dans le disque (étoile et gaz). Le halo possède un paramètre de concentration faible ($c = 5$) comparé aux galaxies de l'Univers local. On affiche la courbe de rotation pour la distribution de masse de chaque composante du modèle. Les échelles de longueur des disques stellaire et gazeux sont respectivement de $r_{stars} = 2.3 \text{ kpc}$ et $r_{gas} = 3.7 \text{ kpc}$, et les deux disques sont tronqués à $r_{cut} = 3r_{gas}$, ce qui affecte la dérivée numérique et fait remonter la courbe de rotation du disque de gaz car cette composante est coupée à un rayon similaire aux étoiles qui ont une longueur d'échelle plus petite.

Cette description simple devrait permettre de créer des modèles de galaxies avec des disques fins en équilibre dynamique. En réalité, une galaxie avec un support dynamique

assuré par de la rotation pure est dynamiquement très instable et développe rapidement des instabilités axisymétriques. Le support dynamique des galaxies est en fait en partie assuré par la dispersion des vitesses de ses étoiles. Ce support dispersif est plus conséquent dans les galaxies elliptiques que dans les galaxies spirales, mais il est impossible de négliger ce support pour créer un modèle dynamiquement stable. En effet, la dispersion des vitesses azimutales permet notamment de protéger les disques en rotation contre ces instabilités axisymétriques qui se développent naturellement. La modélisation numérique de galaxies idéalisées s'appuie donc nécessairement sur les équations de Jeans (Binney et al., 2009), qui permettent de modéliser de manière réaliste le support dispersif des disques. Il s'agit pour chaque particule Lagrangienne d'intégrer ces équations, en supposant que chaque composante cylindrique des vitesses est distribuée suivant une loi Gaussienne, ce qui est une approximation usuelle¹ que j'utilise pour construire mes conditions initiales.

Afin de spécifier entièrement le système, nous supposons que les composantes radiales et verticales de la vitesse locale moyenne sont nulles $\langle v_r \rangle = \langle v_z \rangle = 0$. Cela revient à faire l'hypothèse de la stabilité du modèle de galaxie dans un environnement fermé (i.e. sans apport de masse externe). Les trois composantes du vecteur vitesse de chaque particule Lagrangienne sont déterminées par l'intégration numérique des équations de Jeans, en faisant l'hypothèse que la distribution statistique des vitesses a la forme d'une Gaussienne tri-axiale. Pour chaque particule d'une composante non-collisionnelle du modèle idéalisé, on résout numériquement les équations Jeans :

$$\langle v_z^2 \rangle = \langle v_r^2 \rangle = \frac{1}{\rho} \int_z^\infty \rho(r, z') \frac{\partial \Phi}{\partial z'} dz' , \quad (2.13)$$

avec ρ la fonction de densité volumique de la composante considérée, et $\Phi(r, z)$ le potentiel gravitationnel total du modèle de galaxie idéalisée, i.e. qui tient compte de la masse de toutes les composantes. Pour la composante azimutale, la dispersion des vitesses des particules est spécifiée par l'équation

$$\langle v_\phi^2 \rangle = \langle v_r^2 \rangle + \frac{r}{\rho} \frac{\partial(\rho \langle v_r^2 \rangle)}{\partial r} + r \frac{\partial \Phi}{\partial r} . \quad (2.14)$$

La dispersion des vitesses pour une particule d'une composante située à la coordonnée cylindrique $[r, z]$ peut donc être obtenue grâce à la relation :

$$\sigma_\phi^2 = \langle v_\phi^2 \rangle - \langle v_\phi \rangle^2 , \quad (2.15)$$

$$\sigma_r^2 = \langle v_r^2 \rangle , \quad (2.16)$$

$$\sigma_z^2 = \langle v_z^2 \rangle . \quad (2.17)$$

Le halo de matière noire est généralement modélisé avec un moment angulaire qui n'est pas spécifié par les équations de Jeans. La composante de courant azimutal est fixée de manière à être une petite fraction f_s de la vitesse circulaire, i.e. $\langle v_\phi \rangle = f_s v_c$. La fraction f_s dépend du paramètre de rotation du halo λ ainsi que du paramètre de concentration du halo c (Springel et White, 1999), qui sont les deux paramètres avec le paramètre d'échelle de longueur permettant de caractériser le halo de matière sombre dans un modèle idéalisé de galaxie.

1. Cette hypothèse peut néanmoins ne pas être suffisante dans certaines régions du modèle et peut introduire des instabilités notamment dans les parties centrales du disque et provoquer des effets de relaxation dynamique lors de l'intégration temporelle.

Quant au disque stellaire, il est possible d'utiliser l'approximation de dérive axisymétrique (Binney et al., 2009) qui permet de raccourcir le temps de calcul, bien que cette approximation nécessite de se situer dans le cas d'un disque fin². Cette approximation relie la dispersion de vitesse radiale à la dispersion azimuthale :

$$\sigma_\phi^2 = \frac{\sigma_r^2}{\eta^2}, \quad (2.18)$$

avec

$$\eta^2 = \frac{4}{\kappa^2 r} \frac{\partial \Phi}{\partial r}, \quad (2.19)$$

et

$$\kappa^2 = \frac{3}{r} \frac{\partial \Phi}{\partial r} + \frac{\partial^2 \Phi}{\partial r^2}. \quad (2.20)$$

avec κ la fréquence épicyclique. Dans la figure 2.4, je présente le résultat de ces intégrations numériques des composantes de dispersions de vitesses pour le modèle de galaxie introduit dans la figure 2.3.

La stabilité du disque contre les instabilités axisymétriques est quantifiée par le paramètre de Toomre. Pour le disque stellaire, il s'écrit :

$$Q_{stars} = \frac{\sigma_z \kappa}{3.36 G \Sigma_{stars}}, \quad (2.21)$$

où κ est désigné par le terme de fréquence épicyclique définie dans l'équation 2.20, et Σ_{stars} est la densité surfacique du disque stellaire. Afin de prévenir la formation d'instabilités axisymétriques, on peut définir un paramètre Q_{min} qui peut être utilisé pour définir la stabilité du disque stellaire en imposant une valeur minimale à la dispersion σ_z , de telle manière que $Q_{stars} \geq Q_{min}$. Dans la figure 2.5, je montre la valeur du paramètre de Toomre Q_{stars} dans le plan des coordonnées cylindriques $r-z$ obtenue numériquement pour le modèle de galaxie de la figure 2.3. Le paramètre de Toomre pour le disque de gaz quant à lui s'écrit

$$Q_{gas} = \frac{\sigma_z \kappa}{\pi G \Sigma_{gas}}. \quad (2.22)$$

De manière générale, comme le paramètre de Toomre est proportionnel à σ_z et κ , plonger les disques dans un potentiel gravitationnel sphérique plus prononcé (i.e. augmenter la valeur de c pour le halo de matière sombre) permet d'augmenter Q , et donc de stabiliser le disque.

Finalement, on définit les composantes radiales et verticales de la vitesse pour les particules de gaz du disque avec $v_{r,gas} = v_{z,gas} = 0$ puisque il s'agit d'une composante collisionnelle régie par une équation d'état. La seule composante à spécifier pour le gaz est la vitesse de courant azimuthale, déduite des équations de Euler :

$$\langle v_{\phi,gas} \rangle = r \left(\frac{\partial \Phi}{\partial r} + \frac{1}{\rho_{gas}} \frac{\partial P}{\partial r} \right), \quad (2.23)$$

où P est le terme de pression du gaz.

2. L'approximation de dérive axisymétrique est valide pour des disques relativement fins. Son utilisation dans le cas de disques stellaires plus épais dont une partie du support gravitationnel est assurée par la dispersion de vitesses peut générer des artefacts de relaxation lors de l'évolution du modèle qui peut rendre les conditions initiales instables.

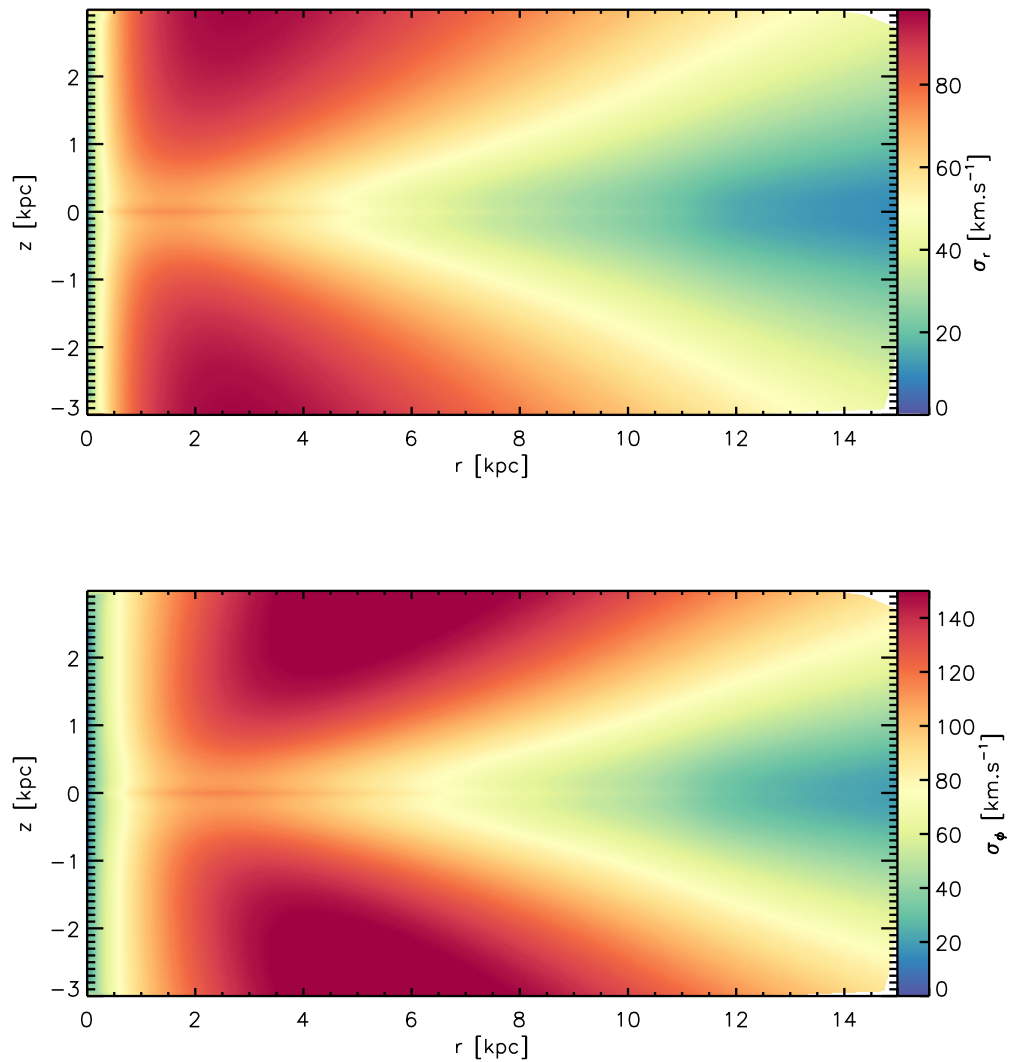


Figure 2.4 – Illustration de la dispersion des vitesses obtenue par l'intégration numérique des équations de Jeans pour les particules stellaires du modèle présenté dans la figure 2.3. En haut : dispersion des vitesses $\sigma_r = \sigma_z$ dans le plan $r - z$ du système de coordonnées cylindriques dont l'axe z est perpendiculaire au plan du disque. En bas : dispersion des vitesses σ_ϕ dans le même plan $r - z$.

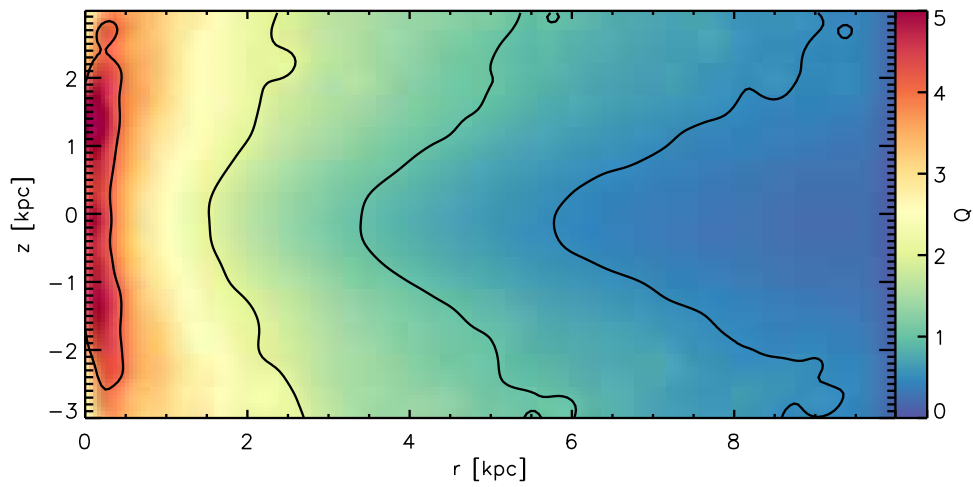


Figure 2.5 – Distribution du paramètre de Toomre Q_{stars} (équation 2.21) pour le disque stellaire présenté dans la figure 2.3 dans le plan des composantes cylindriques $r - z$. Le diagramme est obtenu par l'intégration numérique des équations 2.13, 2.14, et 2.20. Le diagramme ainsi obtenu est lissé par un noyau Gaussien afin de lisser le bruit numérique inhérent aux dérivations numériques du second d'ordre de l'équation 2.20. Les contours délimitent les niveaux $Q_{stars} = [0.5, 1, 2, 4]$. La dispersion des vitesses permet dans les régions où $Q_{stars} > 1$ de prévenir la formation d'instabilités axisymétriques. Afin de prévenir la formation de telles instabilités dans les régions où $Q_{stars} < 1$, il est possible de fixer une valeur minimale pour $\langle v_z^2 \rangle$ afin de s'assurer que $Q_{stars} > Q_{min}$ où Q_{min} est la valeur minimale tolérée pour le paramètre de Toomre.

2.5 Interaction de deux galaxies

Dans cette section, je dérive les équations permettant la mise en place d'une méthode simple pour décrire un système de deux galaxies sur une trajectoire Képlérienne, afin de pouvoir étudier la fusion de galaxies dans le contexte de simulations idéalisées. Notre description de la trajectoire d'un système de deux galaxies se base sur les paramètres suivants :

- m_1 : la masse de la première galaxie.
- m_2 : la masse de la seconde galaxie.
- r_{ini} : la distance initiale entre les deux galaxies.
- r_{peri} : la distance péricentrale, i.e. la distance minimale entre les deux particules Képlériennes.
- e : l'excentricité des trajectoires. Ce paramètre détermine aussi l'énergie orbitale du système $E_{orb} = \frac{v_{ini}^2}{2} \frac{m_1 m_2}{m_1 + m_2} - G \frac{m_1 m_2}{r_{ini}}$.

Le problème gravitationnel à deux corps peut être résolu simplement en utilisant l'approche de la "particule réduite". Dans la formulation du problème, j'ai pris soin d'avoir une écriture valable pour n'importe quelle valeur du paramètre d'excentricité e de la trajectoire Képlérienne. Les paramètres des différentes trajectoires Képlériennes, qui sont en fait assimilables à des sections coniques, sont illustrés dans la figure 2.6 pour plus de clarté. Il est possible de résoudre le problème gravitationnel à deux corps en se plaçant dans le cas équivalent du problème à un corps pour une particule fictive ayant une masse réduite

$$\mu = \frac{m_1 m_2}{m_1 + m_2}, \quad (2.24)$$

et qui orbite autour du centre de masse du système à deux corps à une distance r qui est la distance entre les deux corps. Nous pouvons à présent définir les coefficients de proportionnalités :

$$k_1 = \frac{\mu}{m_1} \quad k_2 = \frac{\mu}{m_2} \quad (2.25)$$

Ces coefficients permettent de passer d'une formulation utilisant l'approche de la particule réduite à une formulation décrivant les deux corps Képlériens. De ce fait, la distance radiale au barycentre du système est :

$$r_1 = k_1 \cdot r_{ini} \quad r_2 = k_2 \cdot r_{ini}. \quad (2.26)$$

Nous définissons à présent le paramètre a de la trajectoire Képlérienne qui, dans le cas d'une ellipse, correspond au demi grand axe.

$$\begin{aligned} a_1 &= k_1 r_{per} & a_2 &= k_2 r_{per} & \text{si} & & e &= 1 \\ a_1 &= \frac{k_1 r_{per}}{|1-e|} & a_2 &= \frac{k_2 r_{per}}{|1-e|} & \text{sinon} & & & \end{aligned} \quad (2.27)$$

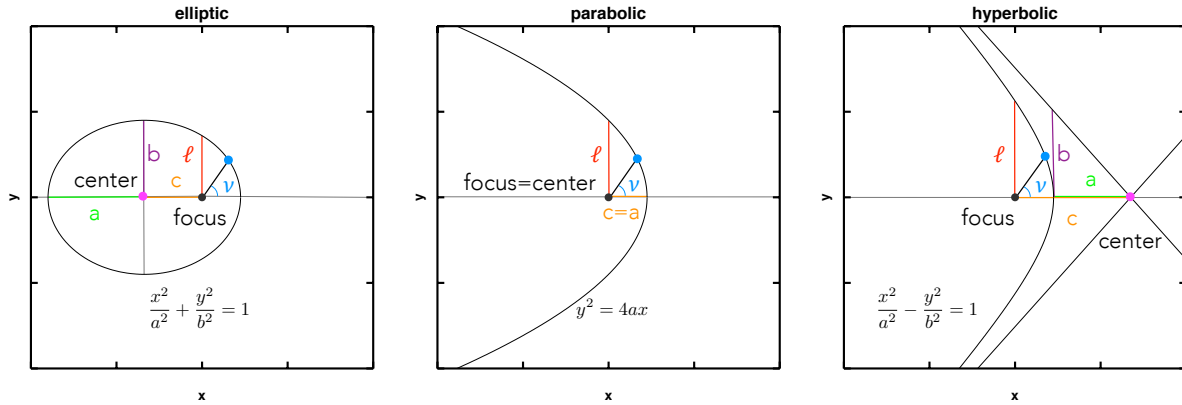


Figure 2.6 – Description des paramètres d'une section conique elliptique (à gauche), parabolique (au milieu) et hyperbolique (à droite). Le demi latus rectum est indiqué en rouge, l'excentricité linéaire est indiquée en orange, le demi grand axe est indiqué en vert, et l'anomalie vraie du point sur la trajectoire Képlérienne est indiquée en bleu. On indique dans chaque cas l'équation Cartésienne de la courbe Képlérienne. Dans le cas hyperbolique, nous traçons les deux asymptotes vers lesquelles convergent les deux branches de l'hyperbole.

Nous considérons à présent le demi latus rectum, qui est le segment passant par le foyer de la section conique considérée, et qui est parallèle à la directrice de la section. Le segment est délimité par la courbe de la-dite conique. Pour chacune des deux galaxies, le demi latus rectum est défini respectivement par :

$$\begin{aligned} \ell_1 = 2a_1 \quad \ell_2 = 2a_2 \quad \text{si} \quad e = 1, \\ \ell_1 = a_1|1 - e^2| \quad \ell_2 = a_2|1 - e^2| \quad \text{sinon.} \end{aligned} \quad (2.28)$$

Avec ces définitions, nous sommes maintenant capables de calculer pour chaque galaxie i la valeur de l'anomalie vraie ν_i :

$$\nu_1 = \arccos\left(\frac{\frac{\ell_1}{r_1} - 1}{e}\right) \quad \nu_2 = \arccos\left(\frac{\frac{\ell_2}{r_2} - 1}{e}\right). \quad (2.29)$$

Comme ce calcul a pour but d'être incorporé dans une méthode globale de création de conditions initiales de galaxies, j'utilise le système de coordonnées cartésiennes qui est la plupart du temps utilisé dans les programmes de simulations numériques. Il s'agit donc du système le plus approprié pour définir la position et la vitesse de chacune des deux galaxies. Dans le plan orbital d'interaction, nous définissons les coordonnées Cartésiennes respectives des deux galaxies réduites à des particules Képlériennes comme suit :

$$\begin{aligned} x_1 = r_1 \cos(\nu_1) \quad y_1 = r_1 \sin(\nu_1), \\ x_2 = -r_2 \cos(\nu_2) \quad y_2 = -r_2 \sin(\nu_2). \end{aligned} \quad (2.30)$$

À noter que le choix du signe devant les coordonnées de la seconde galaxie est complètement arbitraire. Dans cette configuration, les directrices des deux trajectoires Képlériennes

sont parallèles à l'axe y du système de coordonnées Cartésiennes. En passant à la formulation à deux corps du problème, les solutions admissibles pour les orbites des deux particules sont notamment les sections coniques. En utilisant la seconde loi de Kepler, nous trouvons pour la particule réduite :

$$r = \frac{\ell}{1 + e \cos(\nu)}, \quad (2.31)$$

avec r la distance radiale au barycentre de la particule, ℓ le demi latus rectum de la section conique considérée et ν l'anomalie vraie. Le demi latus rectum de la particule réduite est déterminé lors du passage péricentral pour lequel $r = r_{peri}$ et $\nu = 0$.

On peut relier le demi latus rectum avec l'excentricité linéaire $c = ae$ de la trajectoire de la particule réduite (voir figure 2.6) :

$$\ell = \frac{\mu c^2}{Gm_1 m_2}. \quad (2.32)$$

Nous définissons ensuite le paramètre gravitationnel standard :

$$\gamma = G(m_1 + m_2) \quad (2.33)$$

avec G la constante gravitationnelle.

Finalement, les vitesses cartésiennes v_x, v_y des deux galaxies dans le plan orbital d'interaction sont définies par les équations :

$$\begin{aligned} v_{x,1} &= k_1 \sqrt{\frac{\gamma}{\ell}} \sin(\nu_1), & v_{y,1} &= -k_1 \sqrt{\frac{\gamma}{\ell}} [e + \cos(\nu_1)], \\ v_{x,2} &= -k_2 \sqrt{\frac{\gamma}{\ell}} \sin(\nu_2), & v_{y,2} &= k_2 \sqrt{\frac{\gamma}{\ell}} [e + \cos(\nu_2)]. \end{aligned} \quad (2.34)$$

Avec ces définitions, il est possible de définir des trajectoires Képlériennes de n'importe quelle excentricité allant de la trajectoire elliptique à la trajectoire hyperbolique. La figure 2.7 illustre le calcul de trajectoires Képlériennes.

2.6 Implémentation numérique

Les équations décrites ci-dessus permettent de construire plusieurs disques dynamiquement stables, et de les positionner sur une orbite Képlérienne afin d'étudier les propriétés d'une fusion de deux galaxies. L'intégration numérique de ces équations est implémentée dans le code *Disk Initial Conditions Environment* (DICE). L'architecture du code est basée sur le code *starscream*³ qui m'a servi de point de départ pour développer ma propre implémentation numérique. Le code est écrit en C, et est librement disponible sous licence GPL3 à l'adresse <https://code.google.com/p/dice-project/>. La stabilité des conditions initiales pour un disque de type Voie Lactée est illustrée dans la figure 2.8. Le fonctionnement du programme est expliqué en annexe C.

3. <http://code.google.com/p/starscream>

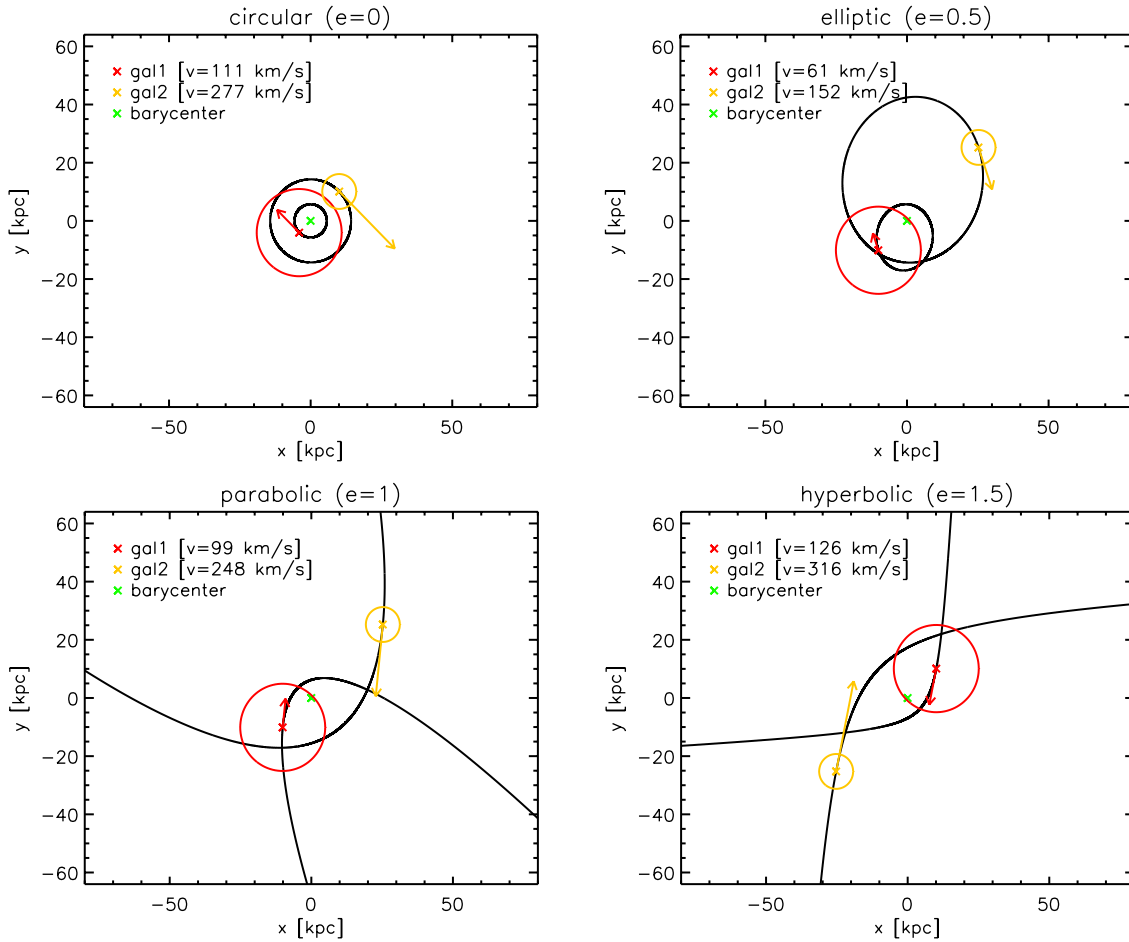


Figure 2.7 – Trajectoires Képlériennes calculées avec diverses excentricités grâce aux systèmes d'équations posés dans cette section. Les trajectoires sont calculées pour deux galaxies de masses totales respectives $M_1 = 5 \times 10^{11} M_\odot$ et $M_2 = 2 \times 10^{11} M_\odot$ dont les disques sont représentés respectivement par un cercle rouge et un cercle orange dont le rapport des tailles est proportionnel au rapport des masses. Dans le cas d'une trajectoire circulaire ($e=0$), elliptique ($e=0.5$), parabolique ($e=1$) et hyperbolique ($e=1.5$) on affiche la trajectoire Képlérienne vue dans le plan orbital d'interaction pour une distance péricentrale $r_{peri}=20$ kpc et une distance de séparation initiale $r_{ini}=50$ kpc (sauf dans le cas $e=0$ où $r_{ini}=r_{peri}=20$ kpc). La vitesse de chacune des galaxies dans le plan orbital est indiquée dans la légende et est représentée par une flèche sur les graphiques. Le barycentre du système est représenté par une croix verte et se situe à $[x=0,y=0]$.

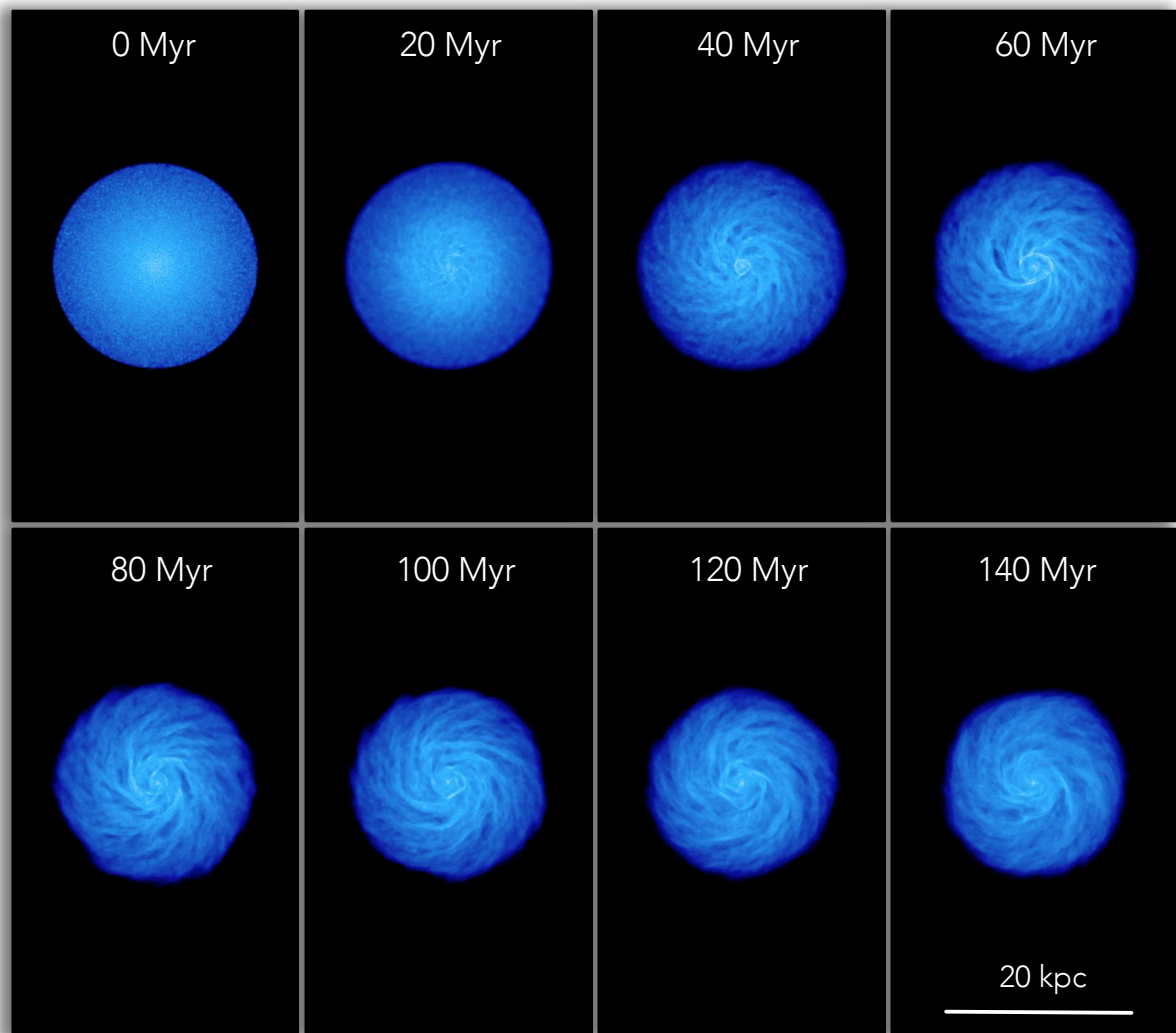


Figure 2.8 – Évolution de la densité du gaz pour une simulation de galaxie comparable à la Voie Lactée vue de face dont les conditions initiales sont produites par DICE. L'intégration temporelle est assurée par le code SPH Gadget2, avec un total de 610 000 particules : 200 000 pour le disque gazeux, 200 000 pour le disque stellaire, 200 000 pour le halo de matière noire, 10 000 pour le bulbe stellaire. La structure du disque gazeux est stable, malgré l'apparition d'un anneau de relaxation inhérent à l'approximation d'une forme Gaussienne pour la distribution des vitesses stellaires.

L'échantillon de simulations MIRAGE



Résumé

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Le projet scientifique majeur de cette thèse a consisté à construire un échantillon de simulations numériques de fusions de galaxies idéalisées. L'objectif initial lors de la définition de cet échantillon était de créer une base de données de simulations qui rendrait possible la détection de signatures cinématiques caractéristiques de la fusion entre des galaxies très riches en gaz, dans différentes configurations orbitales. Dans ce chapitre, je présente l'article qui détaille les étapes de définition, de construction, et d'analyse de l'échantillon Merging and Interacting high-Redshift AMR Galaxies (MIRAGE). Ce chapitre central de cette thèse s'appuie sur les données observationnelles du chapitre 1, sur la méthode de création de modèles de galaxies idéalisées du chapitre 2, et définit les données qui seront utilisées dans les deux chapitres suivants.

Dans le chapitre 1, nous avons vu que les observations spectroscopiques des galaxies "juvéniles" révèlent une fraction importante de systèmes sans support de rotation ordonné. Plus particulièrement, l'échantillon MASSIV que j'ai étudié au cours de cette thèse contient 40% (30/72) de champs de vitesses avec une vitesse maximale de rotation (retrouvée avec la méthode kinemetry) inférieure à 50 km.s^{-1} . Or, les fractions de gaz dans les disques des galaxies de l'Univers jeune ($z > 1$) sont sensiblement plus élevées que dans l'Univers local (Daddi et al., 2010a). Une fraction de gaz élevée est synonyme de formation rapide d'un disque en rotation de par la conservation du moment angulaire. En effet, le gaz étant une composante dissipative, il est capable de rayonner son énergie interne sur des échelles de temps courtes (quelques millions d'années), lui permettant de re-former rapidement une structure adoptant la forme d'un disque en rotation. Les simulations numériques avec fractions de gaz élevées montrent qu'il est difficile de détruire le disque de gaz (Springel et Hernquist, 2005) par une fusion galactique, contrairement au cas à faible fraction de gaz où ce dernier est consommé

durant la flambée de formation stellaire associée à la fusion. Comment expliquer alors autant de systèmes sans support dynamique de rotation ? La qualité des données et le rapport signal-sur-bruit suffisant d'une importante fraction de l'échantillon nous conduit à chercher des explications physiques pour expliquer ce comportement dynamique inattendu. Parmi les différentes hypothèses permettant d'expliquer la nature de ces galaxies, on peut lister :

- Des objets produits d'une fusion récente avec une composante gazeuse non-relaxée.
- Des disques vus dans le plan du ciel de telle manière que seuls les mouvements perpendiculaires au disque soient visibles. Cependant, une distribution aléatoire des disques, ce qui semble confirmé par les simulations cosmologiques, induit une faible probabilité d'observer des galaxies vues de face.
- Une structure sphéroïdale gazeuse dont le support gravitationnel est essentiellement dispersif. Néanmoins, les fortes fractions de gaz conduisent systématiquement à la création d'un disque fin en rotation.
- Une accrétion continue de gaz du milieu circum-galactique avec un taux élevé, qui perturberait le support dynamique du disque.
- Un disque à la morphologie complexe dominée par quelques granules éloignés du centre de masse concentrant la majeure partie de l'émission du gaz ionisé.

Durant les premiers mois de cette thèse, j'ai initialement abordé cette problématique en planifiant une étude des signatures cinématiques de fusion dans l'échantillon de simulations idéalisées GalMer Chilingarian et al. (2010). La base de données GalMer, disponible en ligne¹, compile plusieurs centaines de simulations idéalisées de fusions majeures de galaxies de la séquence de Hubble en étudiant notamment les paramètres orbitaux. Le grand nombre de simulations de la base de donnée GalMer est l'un de ses points forts, et permet des analyses statistiques sur l'évolution d'observables durant la fusion (e.g. Di Matteo et al., 2007, 2008). Bien que l'exploration systématique des orientations des disques se prête particulièrement à l'analyse des signatures de fusions sur la cinématique du gaz, l'étude comparative avec les données observationnelles à grand décalage spectral s'est révélée difficile.

En effet, les fractions de gaz élevées des objets observés peuvent conduire à des instabilités gravitationnelles qui forment des structures dont il est difficile de trouver un équivalent dans l'Univers proche. Cette fragmentation engendre des modifications substantielles dans la manière dont le gaz est distribué durant la coalescence (Bournaud et al., 2011). On observe donc dans les simulations numériques idéalisées de fusions de galaxies deux processus en compétition qui tendent à la fois à reformer le disque après la coalescence, mais aussi à détruire ce dernier par instabilités gravitationnelles. Sur de grandes échelles spatiales, le rayonnement de l'énergie interne du gaz crée un disque fin essentiellement supporté par la rotation. La dissipation de cette énergie interne ne s'arrête pas au moment où le gaz s'écoule dans un disque, le refroidissement se poursuit et fragmente le disque. Ces processus physiques ne sont pas pleinement décrits par les simulations GalMer dont l'objectif n'était d'ailleurs pas de simuler des interactions entre galaxies à grand décalage spectral du fait (i) de la faible fraction de gaz (entre 5 et 20%), (ii) du traitement isotherme du solveur hydrodynamique qui empêche de capturer la condensation et la fragmentation du gaz (iii)

1. <http://galmer.obspm.fr>

et enfin de la faible résolution numérique qui ne permet pas de résoudre la structure verticale du disque, ce qui est indispensable pour s'assurer que la fragmentation observée n'est pas numérique mais physique. C'est pour ces raisons que j'ai fait le choix de construire un nouvel échantillon de simulations de fusion de galaxies riches en gaz, avec un traitement hydrodynamique capable de capturer la fragmentation et la reformation du disque.

De manière naturelle, je me suis orienté vers l'utilisation des simulations idéalisées. Il s'agit de la méthode la plus efficace pour atteindre des résolutions physiques très élevées (on atteint à ce jour des résolutions sub-pc, e.g. Bournaud et al., 2010; Hopkins et al., 2013a; Renaud et al., 2013). Le prix à payer pour atteindre ces résolutions est de se couper d'un environnement cosmologique réaliste durant l'évolution de la simulation, les conditions initiales étant par contre définies par l'environnement cosmologique. Le projet de création d'un échantillon de simulations de fusions riches en gaz s'est initialement construit autour de tests sur des conditions initiales idéalisées. Le contrôle que m'a fourni le développement de DICE sur les modèles de disques m'a permis de définir avec plus de pertinence mon échantillon que si j'avais eu à utiliser des conditions initiales définies pour un projet scientifique différent. J'ai tout d'abord réalisé de nombreux tests permettant de quantifier la stabilité des conditions initiales, ce qui s'est révélé une expérience extrêmement enrichissante. En effet, cela a requis la compréhension de nombreuses méthodes numériques, mais aussi de nombreux paramètres physiques définissant la structure et la stabilité dynamique des galaxies telles qu'on sait les modéliser. Le choix de simuler des galaxies avec des conditions initiales typiques de l'Univers à $z \sim 2$ a nécessité la construction de halos de matière sombre dont les paramètres de concentration sont relativement faibles comparés à ceux de l'Univers local. J'ai choisi d'utiliser $c = 5$ pour tous mes modèles, une valeur en accord avec les mesures dans les simulations cosmologiques (voir figure 3.1). Comme indiqué dans le chapitre 2, un halo peu massif implique un paramètre de Toomre faible et donc un disque par nature instable. Il a donc fallu adapter les stratégies d'initialisation et de relaxation des disques pour contrôler au mieux la stabilité des simulations de disques

Dans cette démarche scientifique, j'ai choisi de favoriser la résolution numérique ainsi que l'ajout de traitements de processus physiques dans le code hydrodynamique `Ramses` aux dépens de l'exploration statistique des paramètres orbitaux de rencontre des galaxies. Le choix de 5 configurations de masse explorant 4 orientations initiales de disques m'a semblé optimal dans l'objectif d'explorer la réponse du gaz pendant une fusion entre deux galaxies. Les contraintes en termes de ressources de temps de calcul m'ont conduit à devoir faire des choix sur les trajectoires des galaxies permettant de se placer dans le cas de durées d'interaction relativement courtes (moins de 1 Gyr). Pour ce faire, j'ai fixé l'énergie orbitale spécifique à une valeur constante pour toutes les configurations de masse ($E^* = -2.85 \times 10^4 \text{ km}^2 \cdot \text{s}^{-2}$) de manière à réduire le nombre de paramètres libres. En intégrant les équations du mouvement, j'ai pu calculer et fixer le temps péricentral, c'est-à-dire le temps nécessaire aux galaxies pour atteindre simultanément le périastre de leur trajectoire Képlérienne, à $t_{\text{peri}} = 250 \text{ Myr}$. Ce temps est suffisamment long pour permettre aux disques de se relaxer, et suffisamment court pour ne pas avoir à dépenser une quantité importante de temps de calcul dans les étapes de pré-fusion. Les configurations orbitales ainsi obtenues sont sensiblement moins énergétiques (i.e. moins excentriques) que celles des simulations cosmologiques N-corps Khochfar et Burkert (2006), mais la distance initiale entre les galaxies, inférieure à 70 kpc dans toutes les simulations MIRAGE, permet de faire l'hypothèse que la friction dynamique a déjà commencé à agir et transférer du moment orbital vers l'énergie interne des halos de matière sombre. Les trajectoires Képlériennes dans le plan orbital d'interaction qui résultent de cette stratégie sont présentées dans la figure 3.2.

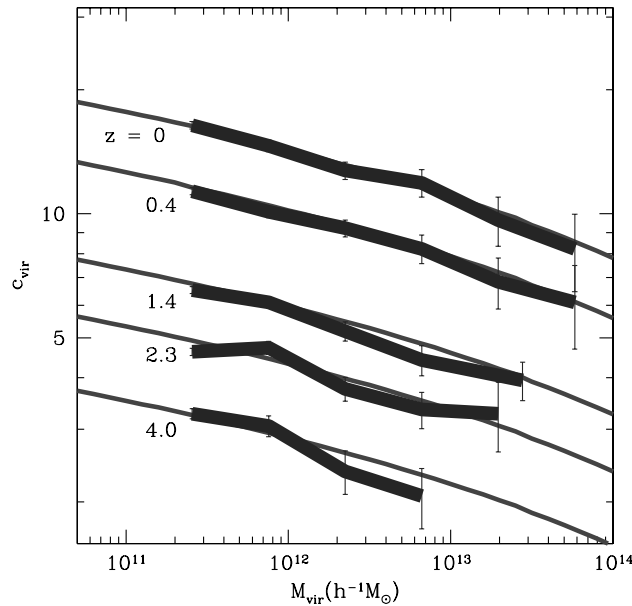


Figure 3.1 – Valeurs médianes du paramètre de concentration c des halos de matière sombre en fonction de la masse de Viriel, pour une simulation cosmologique N-corps de Bullock et al. (2001). Chaque courbe est associée à une mesure pour un décalage spectral donné, et les barres d'erreurs sont calculées pour une statistique de Poisson. Les courbes fines sont les prédictions du modèle étudié dans cette étude. Crédits : Bullock et al. (2001).

L'article ci-dessous présente la définition, la construction et l'analyse de l'échantillon MIRAGE constitué de 20 simulations de fusions de galaxies construites à partir de trois modèles de galaxies ayant des masses baryoniques respectives de 1.4 , 2.5 et $8.8 \times 10^{10} M_{\odot}$ et une fraction de gaz $f_g = 60\%$. Les simulations MIRAGE explorent notamment les orientations initiales des disques, avec des configurations de moment angulaire significativement différentes afin de délimiter au mieux les possibles signatures de fusions de galaxies jeunes dans la cinématique du gaz rémanent. L'évolution des conditions initiales qui sont générées avec DICE est assurée par le code à grille adaptative `Ramses`² avec une résolution spatiale de ~ 7 parsecs dans les cellules les plus raffinées de la grille Cartésienne adaptative. L'implémentation de `Ramses` utilisée pour l'échantillon MIRAGE inclut la formation stellaire, le refroidissement du gaz par émission spontanée des atomes métalliques, l'advection de métallicité dans le solveur hydrodynamique, et un modèle de rétro-action stellaire pour les supernovae, ainsi que le chauffage par photo-ionisation et la pression radiative des étoiles OB (Renaud et al., 2013).

Une moitié des simulations a été réalisée sur le super-calculateur français CURIE³, tandis que la seconde moitié a été réalisée sur le super-calculateur allemand SuperMUC⁴. L'intégration numérique de chacune des simulations a été assurée par 512 processeurs, avec un minimum de 2GB de mémoire vive par processeur. Les simulations modélisant les galaxies les plus massives nécessitent typiquement 20 jours complets de calcul, tandis que les simulations des moins massives requièrent en moyenne 10 jours. Au total, le projet MIRAGE a mobilisé 3.5 millions d'heures de calcul CPU. Toutes les figures en annexe de l'article présenté ci-dessous peuvent être trouvées dans l'annexe A.2 de ce manuscrit.

2. <http://bitbucket.org/rteyssie/ramses>

3. <http://www-hpc.cea.fr/en/complexe/tgcc.htm>

4. <http://www.lrz.de/services/compute/supermuc/>

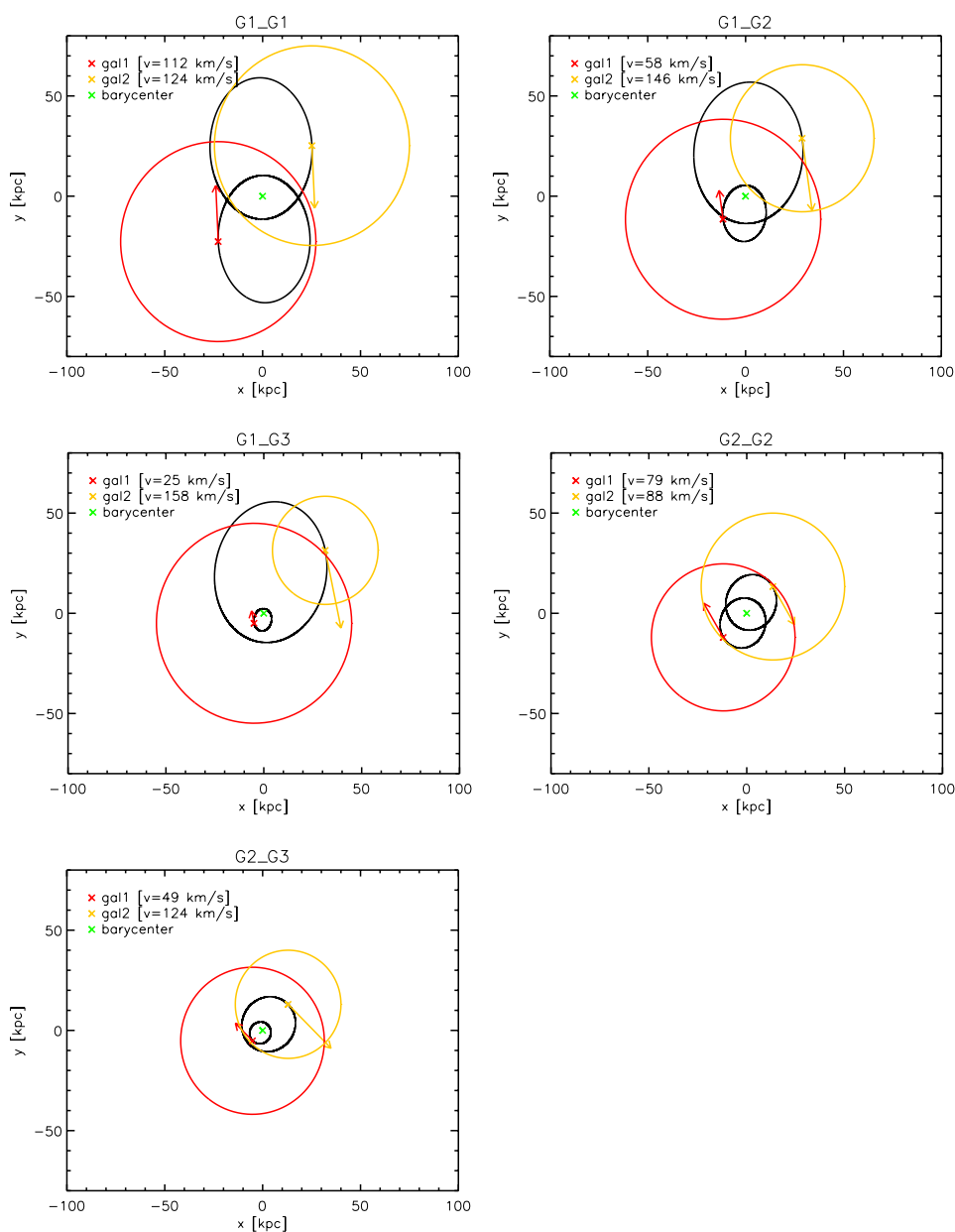


Figure 3.2 – Représentation des orbites Képlériennes initiales des simulations de fusion de l'échantillon MIRAGE dans le plan orbital d'interaction. Les trajectoires Képlériennes de chacun des deux modèles galactiques sont tracées par les ellipses noires. Les cercles rouges et oranges tracent respectivement le rayon de coupure du halo de matière noire pour la première et la deuxième galaxie du système. La vitesse absolue (tangentielle à l'orbite) de chacune des galaxies dans le référentiel Cartésien Galiléen centré sur le barycentre du système est indiquée dans la légende de chaque sous-figure. Dans chaque panneau, le barycentre du système, situé aux coordonnées cartésiennes $[0,0]$, est représenté par une croix verte.

Evolution of the mass, size, and star formation rate in high redshift merging galaxies

MIRAGE - A new sample of simulations with detailed stellar feedback

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ABSTRACT

Context. In Λ -CDM models, galaxies are thought to grow both through continuous cold gas accretion coming from the cosmic web and episodic merger events. The relative importance of these different mechanisms at different cosmic epochs is nevertheless not yet understood well.

Aims. We aim to address questions related to galaxy mass assembly through major and minor wet merging processes in the redshift range $1 < z < 2$, an epoch that corresponds to the peak of cosmic star formation history. A significant fraction of Milky Way-like galaxies are thought to have undergone an unstable clumpy phase at this early stage. We focus on the behavior of the young clumpy disks when galaxies are undergoing gas-rich galaxy mergers.

Methods. Using the adaptive mesh-refinement code RAMSES, we build the Merging and Isolated high redshift Adaptive mesh refinement Galaxies (MIRAGE) sample. It is composed of 20 mergers and 3 isolated idealized disks simulations, which sample disk orientations and merger masses. Our simulations can reach a physical resolution of 7 parsecs, and include star formation, metal line cooling, metallicity advection, and a recent physically-motivated implementation of stellar feedback that encompasses OB-type stars radiative pressure, photo-ionization heating, and supernovae.

Results. The star formation history of isolated disks shows a stochastic star formation rate, which proceeds from the complex behavior of the giant clumps. Our minor and major gas-rich merger simulations do not trigger starbursts, suggesting a saturation of the star formation due to the detailed accounting of stellar feedback processes in a turbulent and clumpy interstellar medium fed by substantial accretion from the circumgalactic medium. Our simulations are close to the normal regime of the disk-like star formation on a Schmidt-Kennicutt diagram. The mass-size relation and its rate of evolution in the redshift range $1 < z < 2$ matches observations, suggesting that the inside-out growth mechanisms of the stellar disk do not necessarily require cold accretion.

Key words. galaxies: evolution – galaxies: formation – galaxies: high redshift – galaxies: star formation – galaxies: interactions – methods: numerical

1. Introduction

Lambda-CDM cosmological simulations tend to show that a major merger is at work shaping galaxy properties at high redshifts (Stewart et al. 2009). Although it is often set as a competitor of the smooth cold gas accretion along cosmic filaments, which is believed to be very efficient at feeding star formation (Dekel et al. 2009a; Kereš et al. 2009b), mergers still contribute to around a third of the baryonic mass assembly history (Brooks et al. 2009; Dekel et al. 2009b). The pioneering work of Toomre & Toomre (1972) first highlighted that disk galaxy mergers are able to drive large amounts of baryons in tidal tails. Mihos & Hernquist (1994) showed that the redistribution of gas can fuel star formation enhancement in the core of the remnants. Furthermore, stars during a merger event are gravitationally heated and can form spheroids (Barnes & Hernquist 1996; Mihos & Hernquist 1996), underlining a convincing link between the late-type and early-type galaxies of the Hubble sequence. Boxy slowly rotating ellipticals, however, probably formed at much higher redshifts through multiple minor mergers or in-situ star formation (Oser et al. 2010; Feldmann et al. 2010; Johansson et al. 2012).

The study of stars and gas kinematics is a good way to detect signatures of merger in the recent history of galaxies (Barnes 2002; Arribas & Colina 2003; Bois et al. 2011), and allows constraining the role of mergers in galaxy mass assembly. The past decade has seen the first resolved observations of galaxies in the redshift range $0.5 < z < 3$ using integral fields unit spectrographs (IFU) (Yang et al. 2008; Förster Schreiber et al. 2009; Law et al. 2009; Gnerucci et al. 2011; Contini et al. 2012), where a peak is observed in the cosmic star formation history. This peak located around $z \sim 2$ (Hopkins & Beacom 2006; Yang et al. 2008) could arise from intense merger activity, since it is an efficient mechanism for producing starbursts in the local Universe. Shapiro et al. (2008) did a kinematical analysis of the high- z IFU SINS sample to determine the fraction of mergers. To calibrate this analysis, a set of local observations (Chemlin et al. 2006; Daigle et al. 2006; Hernandez et al. 2005), hydrodynamical cosmological simulations (Naab et al. 2007), and toy models (Förster Schreiber et al. 2006) were used. The halos in the Naab et al. (2007) cosmological simulations were selected to host a merger around $z = 2$. Although the baryon accretion his-

tory makes these simulations credible in terms of cosmological mass assembly, the resulting low number of halos could be considered as insufficient at statistically detecting various merger signatures.

The GalMer database (Chilingarian et al. 2010) favors a statistical approach with hundreds of idealized merger simulations, which are probing the orbital configurations. GalMer is relevant for studying such merger signatures at low redshift (Di Matteo et al. 2008); however, the low gas fractions makes the comparison with high redshift galaxies impossible. Additionally, simulating the interstellar medium (ISM) of high redshift galaxies requires correctly resolving the high redshift disk scaleheights, which can otherwise artificially prevent the expected Jeans instabilities. Indeed, it is now commonly accepted that high redshift disks are naturally subject to such instabilities (Elmegreen et al. 2009, 2007). The high gas fractions at $z > 1$ (Daddi et al. 2010a; Tacconi et al. 2010) are strongly suspected of driving violent instabilities that fragment the disks into large star-forming clumps (Bournaud et al. 2008) and generate turbulent velocity dispersions (e.g., Epinat et al. 2012; Tacconi et al. 2008). Therefore, the canonical image of smooth extended tidal tails falling onto the merger remnant cannot be valid in the context of gas-rich interactions (Bournaud et al. 2011).

The ability to form such clumps is important to understand the complex behavior of high redshift galaxies, but it is also essential to prevent the overconsumption of gas expected at these very high gas densities from the classical Schmidt law. To match the Kennicutt-Schmidt (KS) relation (Kennicutt 1998) and to have acceptable gas consumption timescales, an efficient stellar feedback is required to deplete the gas reservoir of the star-forming regions. Indeed, cosmological simulations with weak or no feedback models produce galaxies with too many baryons in the galactic plane (Kereš et al. 2009a) when compared to the abundance-matching techniques (Guo et al. 2010). The constraints on the intergalactic medium (IGM) metal enrichment also imply that baryons entered galaxies at some points and underwent star formation (Aguirre et al. 2001). It has been demonstrated that scaling supernovae stellar winds in cosmological simulations to the inverse of the mass of the host galaxy produces models in reasonable agreement with the local mass function (Oppenheimer et al. 2010). It is therefore essential to constrain the parameters controlling the stellar feedback processes in order to better understand the scenarios of galaxy evolution.

To get insight into the various processes of galaxy mass assembly, such as mergers, the Mass Assembly Survey with SINFONI in VVDS (MASSIV, Contini et al. 2012) aims to probe the kinematical and chemical properties of a significant and representative sample of high redshift ($0.9 < z < 1.8$), star-forming galaxies. Observed with the SINFONI integral-field spectrograph at the VLT and built upon a simple selection function, the MASSIV sample provides a set of 84 representatives of normal star-forming galaxies with star formation rates ranging from 5 to $400 M_{\odot} \text{yr}^{-1}$ in the stellar mass regime $10^9 - 10^{11} M_{\odot}$. Compared to other existing high- z IFU surveys, the main advantages of the MASSIV sample are its representativeness since it is flux-selected from the magnitude-limited VVDS survey (Le Fèvre et al. 2005) and its size, which allows probing different mass and star formation rate (SFR) ranges, while keeping enough statistics in each category. Together with the size of the sample, the spatially-resolved data therefore allows galaxy kinematics and chemical properties to be discussed across the full mass and SFR ranges of the survey to derive robust conclusions for galaxy mass assembly on cosmological timescales. By studying strong kinematic signatures of merging and detecting pairs in the first-epoch

MASSIV, Epinat et al. (2012) have shown that the fraction of interacting galaxies is up to at least one third of the sample and that more than a third of the galaxies are non-rotating objects. In addition, there are more non-rotating objects in mergers than in isolated galaxies. This suggests that a significant number of isolated non-rotating objects could be mergers in a transient state in which the gas is not dynamically stable. Furthermore, based on the whole MASSIV sample, López-Sanjuan et al. (2013) find a gas-rich major merger fraction of $\sim 20\%$ in the redshift range $1 < z < 1.8$ and a gas-rich major merger rate of ~ 0.12 . Quantification of the kinematical signatures of interacting galaxies and mergers and the understanding of the high fraction of non-rotating systems, the existence of inverse metallicity gradient in some disks (Queyrel et al. 2012), and more generally, a comprehensive view of the process of formation of turbulent and clumpy gaseous galaxy disks, has motivated building a set of simulations of merging galaxies in the redshift range probed with MASSIV, i.e. the Merging and Isolated high redshift Adaptive mesh refinement Galaxies (MIRAGE) simulations.

We describe, a set of 20 idealized galaxy mergers and three isolated disks using adaptive mesh refinement (AMR) simulations with a physically motivated implementation of stellar feedback¹. This paper focuses on presenting of the MIRAGE sample, the numerical technique employed, and the physical properties deduced. The analysis is extended in a companion paper Bournaud et al. (2013) that presents a study of the clumps properties in the three isolated disk simulations of the MIRAGE sample. The paper is organized as follows. In section 2, we describe the numerical technique used to build our simulation sample. In section 3, we specifically describe the idealized initial conditions generation. For this purpose we introduce the new public code DICE and summarize the different numerical techniques used to generate stable galaxies models. Section 4 reviews the MIRAGE sample definition of galactic models and orbital parameters. Section 5 describes the global properties of the sample. For each simulation, we present the star formation histories, the disk scalelengths evolution, and their position on the KS relation.

2. Simulations

We ran a set of idealized AMR high redshift galaxy simulations. The sample encompasses 20 major and minor galaxy mergers and three isolated disks, with a high gas fraction ($>50\%$) typical of $1 < z < 2$ galaxies (Daddi et al. 2010a), evolved over 800 Myr. In this work, we choose to balance the available computational time between high-resolution and statistical sampling of the orbital parameters to provide new insight into the galactic mass assembly paradigm.

2.1. Numerical technique

To build our numerical merger sample, we use the AMR code RAMSES (Teyssier 2002). The time integration of the dark matter and the stellar component is performed using a particle-mesh (PM) solver, while the gas component evolution is insured by a second-order Godunov integration scheme. The code has proven its ability to model the complexity of interstellar gas on various galaxies simulations (e.g., Dubois & Teyssier 2008; Teyssier et al. 2013). The computational domain of our simulations is a cube with a side $l_{box}=240$ kpc, and the coarsest level of the AMR

¹ Movies of the simulations of the present paper are available at: http://www.youtube.com/playlist?list=PL_oPMhue14ZSyxcuFiJrUXI-6ej8Q7rv7

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grid is $\ell_{min} = 7$, which corresponds to a cartesian grid with $(2^7)^3$ elements and with a cell size of $\Delta x = 1.88$ kpc. The finest AMR cells reach the level $\ell_{max} = 15$, where the cell size corresponds to $\Delta x = 7.3$ pc. The grid resolution is adapted at each coarse time step between the low refinement ($\ell_{min} = 7$) and the high refinement levels ($\ell_{max} = 15$). Each AMR cell is divided into eight new cells if at least one of the following assertions is true: (i) it contains a gas mass greater than $1.5 \times 10^4 M_\odot$, (ii) it contains more than 25 particles (dark matter or stars), or (iii) the local Jeans length is less than four times the current cell size. This quasi-Lagrangian refinement scheme is comparable to the one introduced in Teyssier et al. (2010) and Bournaud et al. (2010).

The star formation is modeled with a Schmidt law triggered when the density ρ_{gas} overcomes the threshold $\rho_0 = 100 \text{ cm}^{-3}$, with an efficiency $\epsilon_\star = 1\%$:

$$\dot{\rho}_\star = \begin{cases} 0 & \text{if } \rho_{gas} < \rho_0 \\ 0 & \text{if } T > 2 \times 10^5 \text{ K} \\ \epsilon_\star \rho_{gas} / t_{ff} & \text{else,} \end{cases} \quad (1)$$

where $\dot{\rho}_\star$ is the local SFR, $t_{ff} = \sqrt{3\pi/(32G\rho_{gas})}$ is the free-fall time computed at the gas density ρ_{gas} , and T is the temperature of the cell considered. AMR cells with temperature greater than $2 \times 10^5 \text{ K}$ are not allowed to form stars.

The gravitational potential is computed using a PM scheme with a maximum level $\ell_{max,part} = 13$ for the grid, which ensures gravitational softening of at least 29 pc for Lagrangian particles. This choice prevents a low number of dark matter particles per cell, often synonymous with N-body relaxation. We use a thermodynamical model modeling gas cooling provided by the detailed balance between atomic fine structure cooling and UV radiation heating from a standard cosmic radiation background by using tabulated cooling and heating rates from Courty & Alimi (2004). In this model, the gas metallicity acts like a scale factor on the cooling rate.

The gas is forced to stay within a specific area in the density-temperature diagram to prevent multiple numerical artifacts (see Fig. 1):

- In the low IGM density regime ($\rho < 10^{-3} \text{ cm}^{-3}$), we ensure a gravo-thermal equilibrium for the gas by introducing a temperature floor in the halo following a gamma polytrope at the virial temperature $T_{min}(\rho) = 4 \times 10^6 (\rho/10^{-3})^{2/3} \text{ K}$, as in Bournaud et al. (2010).
- For densities between $10^{-3} \text{ cm}^{-3} < \rho < 0.3 \text{ cm}^{-3}$, the temperature floor is isothermal and set to $T_{min}(\rho) = T_{floor}$. At full resolution (i.e. $\ell_{max} = 15$), we have $T_{floor} = 300 \text{ K}$. The densest IGM can reach the $\rho = 10^{-3} \text{ cm}^{-3}$ limit and can condense on the gaseous disk.
- For densities above 0.3 cm^{-3} , we use the temperature floor $T_{min}(\rho) = 300 \times (\rho/0.3)^{-1/2} \text{ K}$; this choice allows us to have a dynamical range in the thermal treatment of the gas up to 30 times colder than the slope of the thermodynamical model used in Teyssier et al. (2010) and Bournaud et al. (2010).
- A density-dependent pressure floor is implemented to ensure that the local Jeans length is resolved at least by $n_{Jeans} = 6$ cells in order to avoid numerical fragmentation, as initially proposed by Truelove et al. (1997). This Jeans polytrope acts like a temperature floor for the very dense gas: it overcomes the cooling regime of the temperature floor starting from $\rho = 2.6 \text{ cm}^{-3}$ when the resolution is maximum, i.e. cells with a size of 7.3 pc. The Jeans polytrope is described by the equation $T_{min}(\rho) = \rho G m_H (l_{box} n_{Jeans} / 2^{\ell_{max}})^2 / (\pi k_B \sqrt{32})$, with m_H the proton mass and k_B the Boltzmann constant.

- We impose a maximum temperature for the gas $T_{max} = 10^7 \text{ K}$. Indeed, the clumps generated by Jeans instabilities typical of gas-rich disks (Bournaud et al. 2008) lead to regions of low density inside the disk, where supernovae can explode. This thermal explosion is thus able to produce sound speed greater than 1000 km.s^{-1} , which affects the time step in the Godunov solver. Setting an upper limit to the temperature is not fully conservative in terms of energy, but our choice of T_{max} ensures a viable time step and a reasonably low energy loss. This issue typical of grid codes is handled in the same way in the recent work of Hayward et al. (2013).

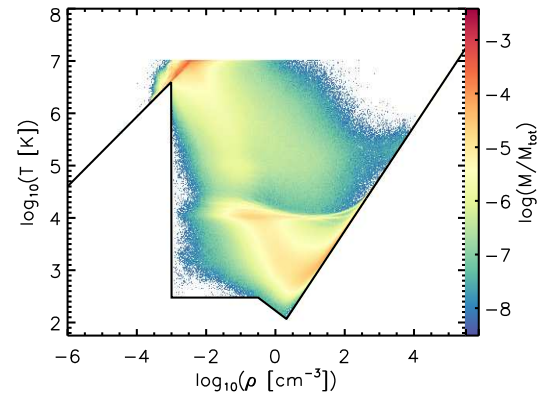


Fig. 1. Density-temperature diagram of the G1 model integrated over 800 Myr (see section 4.1 for a description of the galaxy models). The black line represents the temperature floor described in section 2.1. Each AMR cell contribution to the 2D histogram is weighted by its mass. M represents the gas mass contribution to a bin in the $\rho - T$ plane, color coded on a logscale. M_{tot} is the total gas mass, used as a normalization factor.

Owing to non-periodic boundary conditions applied to the AMR box, we impose a zero density gradient in the hydrodynamical solver at the boundaries. To avoid galaxies passing close to the edges of the box, which could induce numerical artifacts, this gradient is required to have sufficiently large AMR volume to encompass the whole trajectories of both galaxies through the simulation duration.

2.2. Feedback models

Because we do not resolve individual stars, each stellar particle models a population that contains massive OB-type stars with masses $M > 4M_\odot$ (Povich 2012) and which is responsible for injecting energy into the surrounding ISM. Assuming a Salpeter (1955) initial mass function, we consider that a fraction $\eta = 20\%$ of the mass of a stellar particle contributes to stellar feedback, which is effective during 10 Myr after the star particle is spawned. We use the Renaud et al. (2013) physically-motivated model implementation for the OB-type stars feedback, and summarize the three main recipes below.

- Photo-ionization: OB-type stars produce highly energetic photons capable of ionizing the surrounding ISM. Using a simple model for the luminosity of the star, the radius of the Strömgren (1939) sphere is computed according to the mean density of electrons n_e . The gas temperature inside the HII regions is replaced by an isothermal branch at $T_{HII} = 10^4 \text{ K}$.

The radius of the HII sphere is computed via the equation:

$$R_{HII} = \frac{3}{4\pi} \frac{L_*}{n_e^2 \alpha_r}, \quad (2)$$

where L_* is the time-dependent luminosity of the star in terms of ionizing photons, and α_r is the effective recombination rate.

- Radiative pressure: Inside each HII bubble, a kinetic momentum Δv is distributed as a radial velocity kick over the time interval Δt , matching the time step of the coarsest level of the simulation. This velocity kick is computed using ionizing photons momentum, which is considered to be transferred to the gas being ionized, i.e. the gas within the radius R_{HII} of the Strömngren sphere:

$$\Delta v = k \frac{L_* h \nu}{M_{HII} v_c} \Delta t, \quad (3)$$

with h the Planck constant, v_c the speed of light, M_{HII} the gas mass of the bubble affected by the kick, and ν the frequency of the flux representative of the most energetic part of the spectrum of the ionizing source. In this model, it is considered that Lyman- α photons dominate this spectrum, implying $\nu = 2.45 \times 10^{15} \text{ s}^{-1}$. The distribution of the momentum carried by ionizing photons is modeled by the trapping parameter $k = 5$, which basically counts the number of diffusion per ionizing photon and energy loss. This value may appear to be rather high compared to recent works (e.g., Krumholz & Thompson 2012), but is more acceptable once considered that we miss other sources of momentum such as protostellar jets and stellar winds (Dekel & Krumholz 2013).

- Supernova explosions: We follow the implementation of supernova feedback of Dubois & Teyssier (2008): the OB-type star population that reaches 10 Myr transforms into supernovae (SNe) and releases energy, mass, and metals into the nearest gas cell. The gas that surrounds the supernovae receives a fraction $\eta = 20\%$ of the stellar particle mass, as well as a specific energy $E_{SN} = 2 \times 10^{51} \text{ ergs}/10M_\odot$, which is the product of the thermo-nuclear reactions. The energy injected by each SN is higher by a factor of two compared to some other works (e.g., Teyssier et al. 2013; Dubois et al. 2012), but simulations of individual type II SN releasing such energy could be frequent in the early Universe (Joggerst et al. 2010). Moreover, the use of an IMF with a lower statistical contribution of low mass stars would imply higher values for η (e.g. $\eta \simeq 35\%$ for Kroupa 2001 IMF), which balances our choice of a high value for E_{SN} . Each supernovae event also releases into the surrounding ISM metals derived from the nucleosynthesis following the equation

$$Z = y + (1 - y)Z_{ini}, \quad (4)$$

with Z the mass fraction of metals in the gas, Z_{ini} the initial metal fraction of the supernova host, and y the yield that is set to $y = 0.1$, as in Dubois et al. (2012).

To account for non-thermal processes due to gas turbulence on subparsec scales, we follow the revised feedback prescriptions of Teyssier et al. (2013). The numerical implementation is similar to introducing a delayed cooling in the Sedov blast wave solution. At each coarse time step, the fraction of gas released by SNe in AMR cells is evaluated in a passively advected scalar; the gas metal line cooling is switched off as

long as

$$\frac{m_{ejecta}}{m} > 10^{-3}, \quad (5)$$

with m the gas mass of the cell, and m_{ejecta} the total mass of the gas ejected by SNe in the same cell. To model the turbulence dissipation, the mass of gas contributing to the Sedov blast wave is lowered by a factor γ at each coarse time step following

$$\gamma = \exp\left(-\frac{dt_{cool}}{t_{dissip}}\right), \quad (6)$$

with dt_{cool} the cooling time step, and t_{dissip} a typical timescale for the turbulence induced by the detonation. The dissipation timescale for the unresolved subgrid turbulent structures is the crossing time (Mac Low 1999), i.e the ratio of the numerical resolution over the velocity dispersion. Since our simulations are able to resolve structures down to 7 pc, we presume that the non-thermal velocity dispersion in the smallest AMR cells is close to 5 km.s^{-1} , which is a typical supersonic speed in regions of star formation with gas temperatures below 10^3 K (Hennebelle & Falgarone 2012). Under these assumptions, we set $t_{dissip} = 2 \text{ Myr}$.

Our feedback model does not assume the systematic destruction of the clumps by the star formation bursts following their formation, unlike what is done in some other works (e.g., Hopkins et al. 2013; Genel et al. 2012). Smaller clumps are subject to disruption, but larger clumps may survive such thermal energy injection. This model clearly favors the scenario of long-lived star-forming clumps, which we aim to address in this study.

3. DICE: a new environment for building disk initial conditions

The initial conditions of the MIRAGE sample are constructed using software developed for the purpose of the task, named disk initial conditions environment (DICE). DICE is an implementation of the numerical methods described in Springel et al. (2005a). It is able to set up multiple idealized galaxies in a user friendly context. The software is open source and available online².

3.1. Density distributions

DICE initial conditions are generated using Lagrangian particles whose distributions are built using a Metropolis-Hasting Monte-Carlo Markov Chain algorithm (Metropolis et al. 1953). The strength of this algorithm lies in its ability to build a distribution for a sample of Lagrangian particles having only the knowledge of the probability distribution function. After having initialized the first Lagrangian particles of each component (disk, bulge, gas, halo, etc.) to a probable location, the algorithm loops over the desired number of Lagrangian particles and iteratively produces a candidate position for each of them. The probability of setting a Lagrangian particle to the randomly picked candidate cartesian position \mathbf{x}' depends on the cartesian position \mathbf{x} of the previous particle in the loop, and is written as

$$\alpha(\mathbf{x}, \mathbf{x}') = \min\left(1, \frac{\rho(\mathbf{x})P(\mathbf{x}, \mathbf{x}')}{\rho(\mathbf{x}')P(\mathbf{x}', \mathbf{x})}\right), \quad (7)$$

² <http://code.google.com/p/dice-project/>

with $\rho(\mathbf{x})$ the density function of the considered component at the position \mathbf{x} , and $P(\mathbf{x}, \mathbf{x}')$ the probability of placing the particle at \mathbf{x}' considering the position \mathbf{x} of the previous particle. Indeed, our implementation uses a Gaussian walk, meaning that the candidate coordinates are generated using the rule

$$\mathbf{x}' = \mathbf{x} + \sigma W, \quad (8)$$

where W is a standard Gaussian random variable, and σ a dispersion factor tuned to a fixed fraction of the targeted scalelength of the component to build, ensuring satisfying convergence. For each particle, a uniform random value $\tau \in [0, 1]$ is picked, and the position of the Lagrangian particle is set to \mathbf{x}' if $\tau \leq \alpha$ and to \mathbf{x} , otherwise. The first 5% of the iterations to build the distribution are not taken into account because they are considered as a “burning period” to account for any eventual poor choice of initial values.

To fit the system in the finite AMR domain, we cut the density profiles of all the components. We apply these cuts using an exponential truncation profile at the edges of each component, in order to prevent strong discontinuities nearby the cut region, which would make the numerical differentiation quite unstable. The scalelength of the exponential truncation profile is set to be one percent of the gas disk scaleheight.

3.2. Gravitational potential

To set up the velocities in our initial conditions, we compute the gravitational potential using a PM technique. We first interpolate the densities of all the components onto a cartesian grid using a cloud-in-cell scheme. We compute the gravitational potential Φ by solving the Poisson equation:

$$\Phi(\mathbf{x}) = \int \mathcal{G}(\mathbf{x}, \mathbf{x}') 4\pi\rho(\mathbf{x}') d^3\mathbf{x}', \quad (9)$$

where \mathcal{G} is the Green function, and ρ is the density function of all the mass components interpolated on the cartesian grid. We compute this integral by performing a simple product on the Fourier plane, which is equivalent to a convolution in the real plane. We eliminate the periodicity associated to the fast Fourier transform algorithm using the zero-padding technique described in Hockney & Eastwood (1988).

3.3. Velocities

To fully describe our system, we assume that the mean radial and vertical velocities $\langle v_r \rangle$ and $\langle v_z \rangle$ are equal to zero. The velocities of each Lagrangian particle are determined by integrating the Jeans equations (Binney et al. 2009), assuming that the velocity distribution is shaped as a tri-axial Gaussian. For the dark matter halo and the stellar bulge, we numerically solve the equations

$$\langle v_z^2 \rangle = \langle v_r^2 \rangle = \frac{1}{\rho} \int_z^\infty \rho(r, z') \frac{\partial \Phi}{\partial z'} dz', \quad (10)$$

$$\langle v_\phi^2 \rangle = \langle v_r^2 \rangle + \frac{r}{\rho} \frac{\partial (\rho \langle v_r^2 \rangle)}{\partial r} + r \frac{\partial \Phi}{\partial r}. \quad (11)$$

The velocity dispersion can thus be computed using the relation

$$\sigma_\phi^2 = \langle v_\phi^2 \rangle - \langle v_\phi \rangle^2. \quad (12)$$

The dark matter halo is generally described with an angular momentum that is not specified by the Jeans equations. The streaming component is set to be a low fraction f_s of the circular velocity: i.e., $\langle v_\phi \rangle = f_s v_c$. The fraction f_s depends on the halo spin parameter λ and the halo concentration parameter c (Springel & White 1999), which are used as input parameters in our implementation.

For the stellar disk, we choose to use the axisymmetric drift approximation (Binney et al. 2009), which allows fast computation, although we caution against the risk of using this approximation with thick and dispersion supported disks³. This approximation relates the radial Gaussian dispersion to the azimuthal one:

$$\sigma_\phi^2 = \frac{\sigma_r^2}{\eta^2}, \quad (13)$$

with

$$\eta^2 = \frac{4}{r} \frac{\partial \Phi}{\partial r} \left(\frac{3}{r} \frac{\partial \Phi}{\partial r} + \frac{\partial^2 \Phi}{\partial r^2} \right)^{-1}. \quad (14)$$

The Toomre parameter for the stellar disk is written as

$$Q_{stars} = \frac{\sigma_z \kappa}{3.36 G \Sigma_{stars}}, \quad (15)$$

where κ is the so-called epicyclic frequency, and Σ_{stars} is the surface density of the stellar disk. It is used to control the stability of the stellar disk by setting a minimum value for the velocity dispersion σ_z which prevents the local Toomre parameter from going below a given limit of 1.5 in the initial conditions of our simulations, although this parametrization cannot prevent the natural fragmentation of the gaseous disk at later stages.

The only component to specify for the gas is the azimuthal streaming velocity, derived from the Euler equation:

$$\langle v_{\phi, gas} \rangle = r \left(\frac{\partial \Phi}{\partial r} + \frac{1}{\rho_{gas}} \frac{\partial P}{\partial r} \right), \quad (16)$$

where P is the gas pressure.

3.4. Keplerian trajectories

DICE is also able to set up the Keplerian trajectories of two galaxies involved in an encounter. Using the reduced particle approach, we can setup the position of the two galaxies with only three input parameters: (i) the initial distance between the two galaxies r_{ini} ; (ii) the pericentral distance, r_{peri} i.e. the distance between the two galaxies when they reach the periapsis of the Keplerian orbit; (iii) the eccentricity of the trajectories, which are equal for both of the galaxies. The position of the barycenter of each galaxy in the orbital plane can be expressed in cartesian coordinates as

$$\begin{aligned} x_1 &= r_1 \cos(\psi_1), & y_1 &= r_1 \sin(\psi_1), \\ x_2 &= -r_2 \cos(\psi_2), & y_2 &= -r_2 \sin(\psi_2), \end{aligned} \quad (17)$$

³ The axisymmetric drift approximation is valid for relatively thin disks. Using this approximation for thicker disks supported by velocity dispersions might generate relaxation effects that would make the initial conditions unstable.

with ψ_1 and ψ_2 the true anomaly of the first and second galaxy, respectively. The cartesian velocities v_x, v_y of the two galaxies in the orbital plane are computed using

$$\begin{aligned} v_{x,1} &= k_1 \sqrt{\frac{\gamma}{\mathcal{L}}} \sin(\psi_1), & v_{y,1} &= -k_1 \sqrt{\frac{\gamma}{\mathcal{L}}} [e + \cos(\psi_1)], \\ v_{x,2} &= -k_2 \sqrt{\frac{\gamma}{\mathcal{L}}} \sin(\psi_2), & v_{y,2} &= k_2 \sqrt{\frac{\gamma}{\mathcal{L}}} [e + \cos(\psi_2)], \end{aligned} \quad (18)$$

with k_i the mass fraction of the i -galaxy compared to the total mass of the system, γ the standard gravitational parameter, \mathcal{L} the semi-latus rectum of the reduced particle of the system, and e the eccentricity of the orbits. With these definitions, it is possible to set trajectories for any eccentricity. This parametrization holds for point-mass particles, while galaxies are extended objects that undergo dynamical friction. The galaxies quickly deviates from their initial trajectories because of the transfer of orbital energy towards the energy of each galaxy, which can lead to coalescence.

4. Sample definition

4.1. Galaxy models

The different parameters of our disk initial conditions are summarized in Table 1. We set up three idealized galaxy models based on the MASSIV sample stellar mass histogram (Contini et al. 2012). The choice of the initial stellar masses of our simulations was made in order to sample this histogram with all the available snapshots, i.e. in the redshift range $1 < z < 2$. We chose to build our sample out of three disk models with the respective stellar masses: $\log(M_\star/M_\odot) = 9.8$ for our low mass disk, $\log(M_\star/M_\odot) = 10.2$ for our intermediate mass disk, and $\log(M_\star/M_\odot) = 10.6$ for our high mass disk. All of our models have a stellar disk and a gaseous disk with an initial gas fraction $f_g=65\%$. The stellar density profile is written as

$$\rho_{stars}(r, z) = \frac{M_{stars}}{2\pi h_{stars}^2} \exp\left(-\frac{r}{r_{stars}}\right) \exp\left(-\frac{z}{h_{stars}}\right), \quad (19)$$

with r_{stars} the scalelength of the stellar disk, h_{stars} the scaleheight of the stellar disk, and M_{stars} is the uncut stellar disk mass. We use the exact same exponential profile to set up the gaseous disk, with scalelengths 1.68 times shorter than the stellar counterpart as measured in the MASSIV sample data (Vergani et al. 2012). We initialize the metallicity in the gas cells modeling the ISM of the disks following an exponential profile to be consistent with the previous prescriptions:

$$Z(r) = Z_{core} \exp\left(-\frac{r}{r_{metal}}\right). \quad (20)$$

We choose to have negative initial metallicity gradients, with values of r_{metal} equal to the gaseous disk scalelength. The fraction of metals in the center $Z(r=0) = Z_{core}$ of each model is chosen to follow the mass-metallicity relation at $z=2$ found in Erb et al. (2006). Such a choice combined with the exponential profile provides integrated metallicities that are 50 percent lower than the mass-metallicity relation at $z=2$ for starburst galaxies, but this choice is consistent with our aim of modeling normal star-forming galaxies. The numerical implementation of metallicity treatment of the stellar particles ignores the stars present in the initial conditions. It is therefore not required to set a metallicity profile for these stars.

Dark matter halos were modeled using a Hernquist (1990) profile, with a spin parameter set close to the conservative value with $\lambda = 0.05$ (Warren et al. 1992; Mo et al. 1998):

$$\rho_{halo}(r) = \frac{M_{halo}}{2\pi} \frac{a}{r(r+a)^3}, \quad (21)$$

$$a = r_{halo} \sqrt{2 \left(\ln(1+c) - \frac{c}{1+c} \right)}, \quad (22)$$

where M_{halo} is the total dark matter mass, a is the halo scalelength and r_{halo} the scalelength for an equivalent Navarro, Frenk, & White (1997) halo with the same dark matter mass within r_{200} (Springel et al. 2005b). We can therefore define our halo with the frequently used concentration parameter c , which is set to a value $c = 5$ as measured at $z \sim 2$ in N-body cosmological simulations (Bullock et al. 2001). We do not consider the mass dependence of the halo concentration function to ensure that our simulations are comparable in terms of disk instability between each other.

Finally, a bulge enclosing 8% of the total initial stellar mass is modeled again using a Hernquist profile, with a scalelength set to be equal to 20% of the stellar disk scalelength.

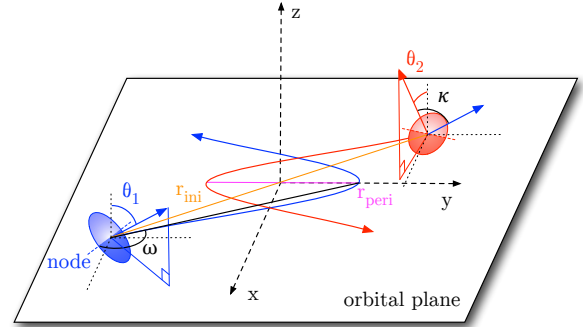


Fig. 2. Orbital geometry used in our simulation sample. Four angles define the geometry of the interaction: θ_1 , θ_2 , κ , and ω . The pericentric argument ω is defined as the angle between the line of nodes (intersection between the orbital plane and the galactic plane) and separation vector at pericenter (black line). The blue/red arrows display the spin orientation for the first/second galaxy. The blue/red curves represent the trajectory of the first/second galaxy in the orbital plane (x,y). The centers of the two galaxies also lie in the orbital plane. The darkest parts of the disks lie under the orbital plane.

4.2. Orbital parameters

The MIRAGE sample is designed to constrain the kinematical signatures induced by a galaxy merger on rotating gas-rich disks. To this purpose, we built a sample that explores probable disks orientations that are likely to produce a wide range of merger kinematical signatures. It has been statistically demonstrated using dark matter cosmological simulations that the spin vectors of the dark matter halos are not correlated one to the other when considering two progenitors as Keplerian particles (Khochfar & Burkert 2006). We use this result to assume that no spin orientation configuration is statistically favored. Our galaxy models are placed on Keplerian orbits using θ_1 the angle between the spin vector of the first galaxy and the orbital plane, θ_2 the angle between the spin vector of the second galaxy and the orbital plane, and κ the angle between the spin vector of the first galaxy and the

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	G1	G2	G3
Virial quantities			
1. $\log_{10}(M_{stars} [M_{\odot}])$	10.60	10.20	9.80
2. R_{200} [kpc]	99.8	73.4	54.0
3. M_{200} [$10^{10} M_{\odot}$]	102.4	40.8	16.2
4. V_{200} [km.s^{-1}]	210.1	154.6	113.7
Scalelengths [kpc]			
5. r_{stars}	2.28	1.62	1.15
6. r_{gas}	3.71	2.64	1.88
7. h_{stars}	0.46	0.32	0.23
8. h_{gas}	0.19	0.13	0.09
9. r_{bulge}	0.46	0.32	0.23
10. r_{halo}	19.95	14.68	10.80
11. $r_{cut,stars}$	11.13	7.92	5.63
12. $r_{cut,gas}$	11.13	1.94	1.38
13. $h_{cut,stars}$	2.73	7.92	5.63
14. $h_{cut,gas}$	0.56	0.40	0.28
15. $r_{cut,bulge}$	2.28	1.62	1.151
16. $r_{cut,halo}$	49.88	36.69	26.99
17. r_{metal}	3.71	2.64	1.88
Mass fractions			
18. f_g	0.65	0.65	0.65
19. f_b	0.10	0.10	0.10
20. m_d	0.10	0.10	0.10
Collision-less particles [10^6]			
21. N_{disk}	2.00	0.80	0.32
22. N_{halo}	2.00	0.80	0.32
23. N_{bulge}	0.22	0.09	0.04
Various quantities			
24. Q_{min}	1.5	1.5	1.5
25. c	5	5	5
26. Z_{core}	0.705	0.599	0.479

Table 1. Physical properties of the three high redshift disk models (G1,G2,G3). All the quantities based on the cosmology use $\Omega_{\Lambda} = 0.7$ and $\Omega_m = 0.3$ and $z=2$.

1. M_{stars} is stellar mass. 2. Virial radius (radius at which the density of the halo reaches 200 times the critical density of the Universe). 3. Cumulated mass at the virial radius. 4. Circular velocity at the virial radius. 5. Stellar disk scalelength. 6. Gaseous disk scalelength. 7. Stellar disk scaleheight. 8. Gas disk scaleheight. 9. Stellar bulge scalelength. 10. Dark matter halo scalelength. 11. Stellar disk radial cut. 12. Stellar disk azimuthal cut. 13. Gas disk radial cut. 14. Gas disk azimuthal cut. 15. Stellar bulge radial cut. 16. Dark matter halo radial cut. 17. Metallicity scalelength. 18. Gas fraction. 19. Stellar bulge mass fraction. 20. Baryonic mass fraction: a mass fraction m_d of M_{200} mass is considered to be in a disk. 21. Number of particles in stellar disk. 22. Number of particles in dark matter halo. 23. Number of particles in stellar bulge. 24. Minimal value for the Toomre stability parameter in the initial conditions. 25. Concentration parameter of the halo. 26. Fraction of metals in the gas at the center of the galaxy in units of solar metallicity.

second one (see Fig. 2). If these angles are uncorrelated, the normalized spin vectors are distributed uniformly over the surface of a sphere. Consequently, all the spin orientations are equally probable. If one considers a random sampling of these disk orientations using a small finite solid angle, having the spin vector coplanar to the orbital plane produces the most configurations. Therefore, we favor configurations where we have at least one spin vector in the orbital plane, i.e. $\theta_1 = 90^\circ$ in all the configurations. We specifically avoid configurations where both of the disks are in the orbital plane because they are highly unlikely and are subject to strong resonances that are not statistically relevant.

We assume that the fourth angle ω that orients the first galaxy with respect to its line of node (Toomre & Toomre 1972) might not affect the kinematics and the shape of the merger remnant since this parameter does not affect the total angular momentum of the system. Consequently, we arbitrarily chose to have the spin vector of the first galaxy always collinear to its Keplerian particle velocity vector. We defined each orbit name with an identifier referring to the angles θ_1 , θ_2 , and κ (see Table 2).

The choice of studying a wide range of spin vector orientations was motivated by the requirement of detecting extreme signatures and binding the kinematical and morphological parameters of the merger remnants. However, we introduced a random angle δ when setting up the spin vector of our galaxies, picked using a uniform distribution introducing a $\pm 5^\circ$ uncertainty. This method was implemented to prevent alignment with the AMR grid, which could produce spurious effects. The slight misalignment also increases the numerical diffusion typical of grid codes, which in our case can help relax our initial conditions.

The pericenter distance, i.e. the distance between the two galactic centers at the time of the closest approach along the Keplerian trajectory, is chosen to be $r_{peri} = r_{1,cut,gas} + r_{2,cut,gas}$, where $r_{1,cut,gas}$ and $r_{2,cut,gas}$ are the cut radii of the first and the second galaxies, respectively. This parametrization ensures that disks do not collide on the first pericentral passage, even if the disks are already fragmented at the pericentral time. This choice is supported by the argument that low pericentral distances are not statistically relevant because the collision cross section is directly proportional to the value of this parameter; i.e., low pericentral distance is less probable. By specifying one value for the specific orbital energy of the system, one can compute the eccentricities of the orbits. To better understand the effects of the interaction parameters on the kinematics of our merger remnants, we set this parameter to a fixed negative value :

$$E^* = \frac{v_{ini}^2}{2} - \frac{G(m_1 + m_2)}{r_{ini}} = -2.85 \times 10^4 \text{ km}^2 \cdot \text{s}^{-2}, \quad (23)$$

where E^* is the specific orbital energy, v_{ini} the initial relative velocity of the galaxies, and r_{ini} the initial distance between the galaxies. This negative specific orbital energy means that all of our trajectories are elliptic ($e < 1$). The parameters of the Keplerian orbits are listed in Table 3. We acknowledge that such low eccentricities might not be statistically relevant (Khochfar & Burkert 2006), but we do save computational time.

orbit label	θ_1	θ_2	κ
90_90_90	90°	90°	90°
90_90_0	90°	90°	0°
90_90_180	90°	90°	180°
90_0_90	90°	0°	90°

Table 2. Orbital angles describing the four orbits studied in this paper. We introduce a random deviation $|\delta| < 5^\circ$ (not given in the table) in our merger setup to avoid over symmetry of our simulations. The orbit name is the concatenation of the angles θ_1 , θ_2 , and κ .

Finally, we define the initial distance between galaxies with a conservative expression through different merger masses, using the pericenter time, i.e., the time for the galaxies to reach the pericenter with $t_{peri}=250$ Myr. The choice of the specific orbital energy is achieved in order to be able to set t_{peri} to 250 Myr with elliptic orbits $e < 1$. Because the dynamical times of all the models are close, this formulation ensures that the models relax synchronously before the start of the interaction (see

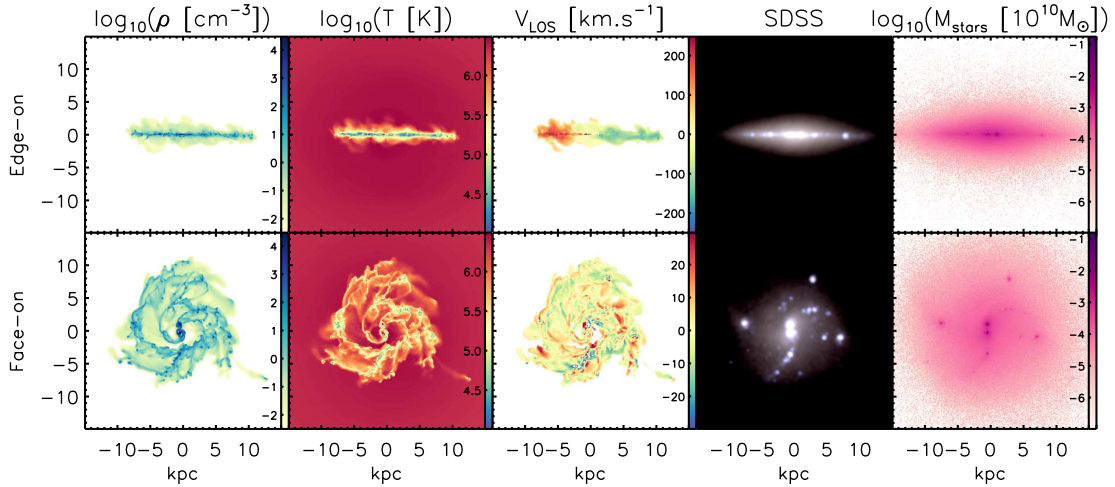


Fig. 3. Maps for the G2 model after 400 Myr of evolution. From left to right: mass-weighted mean gas density, mass-weighted mean gas temperature, mass-weighted mean gas radial velocity, SDSS u/g/r mock observation built from the STARBURST99 model using stellar particles age and mass and assuming solar metallicity, and stellar mass map. The upper line presents an edge-on view, while the bottom line displays a face-on view.

	r_{ini} [kpc]	v_{ini} [km.s ⁻¹]	r_{peri} [kpc]	e	E [10 ⁴ kg.km ² .s ⁻²]
G1_G1	68	237	21.8	0.67	-63.1
G1_G2	57	204	19.1	0.61	-38.0
G1_G3	52	183	16.8	0.59	-18.2
G2_G3	26	173	13.6	0.31	-15.1
G2_G2	36	167	15.5	0.42	-25.1

Table 3. Orbital parameters of the five configurations explored in the MIRAGE sample. These parameters are obtained using $E^* = -2.85 \times 10^4 \text{ km}^2.\text{s}^{-2}$ and $t_{peri}=250 \text{ Myr}$. r_{ini} is the initial distance between the two galaxies, v_{ini} is the initial relative velocity of the two galaxies, r_{peri} is the pericenter distance, e is the eccentricity of the orbits, and E is the orbital energy of the system.

section 5.1). Our sample encompasses 20 merger configurations (four sets of orbital angles, five sets of orbital parameters due to different galaxy masses), to which we add the three isolated disk models to have a reference for secular evolution. We have excluded the G3_G3 interaction to save computation time, since the relative resolution on the merger remnant is coarser than any other cases.

4.3. Environment

We aim to simulate the accretion from an idealized hot gaseous halo surrounding the galactic disks. To this purpose, we model the intergalactic medium (IGM) by setting an initial minimum gas density $\rho_{IGM} = 2.3 \times 10^{-4} \text{ cm}^{-3}$ within the AMR box. The gas present in the IGM is initialized with no velocity, so that it collapses towards the central potential well at the free-fall velocity. After a dynamical time, the gas halo reaches a state close to a spherical hydrostatic equilibrium where the densest regions are allowed to cool down. The zero gradient condition imposed in the grid boundaries implies a continuous injection of pristine gas on the boundaries of the AMR box.

5. Global evolution of physical properties

Figure 3 shows the morphology of the gas and the stars after 400 Myr of evolution along two orthogonal line-of-sight (LOS) for the simulation G2. With the first LOS, we see the disk edge-on, while the second LOS provides a face-on view. For each LOS, the gas density, temperature, the morphology of the stellar component through a rest-frame SDSS mock composite (*ugr* bands), and the stellar mass maps are displayed. We used a pixel of $0.396''$, and we projected our simulations to a luminous distance of 45 Mpc, which gives a pixel size of 0.12 kpc assuming WMAP9 cosmological parameters values. The physical quantities computed for the gas are all mass-weighted averages along the LOS. The stellar emission is computed using the STARBURST99 model (Leitherer et al. 1999) given the age and the mass of each particle. Unlike Hopkins et al. (2013), we chose to neglect the dust absorption in the building of the SDSS mock images to emphasize the stellar light distribution. Projections misaligned with the AMR grid are always difficult to build. To palliate this common issue, we used multiple convolutions with smoothing kernel sizes adapted to the cell sizes.

The projections in Fig. 3 show a disk with clumps lying in a turbulent medium. The most massive clumps reach masses of $\sim 10^9 M_{\odot}$ (Bournaud et al. 2013). We observe there a gaseous disk thickened by stellar feedback. The edge-on velocity field nevertheless shows clear ordered rotation. The clumps concentrate most of the stellar emission due to young stars, since they host most of the star formation. Figure 4 further emphasizes this highly complex behavior of the gas with substantial turbulence and disk instabilities by showing the mass-weighted average density of one of the most massive merger simulation in our sample (G1_G1_90_90_0). We observe star-forming clumps wandering in a very turbulent ISM where the spiral structures are continuously destroyed by the cooling induced fragmentation and the thermal energy injection from stellar feedback. The edge-on view displays a disk thickened by the tidal torque induced by the recent merger. In the appendices of this paper, projections similar to Fig. 3 are given for three simulations of

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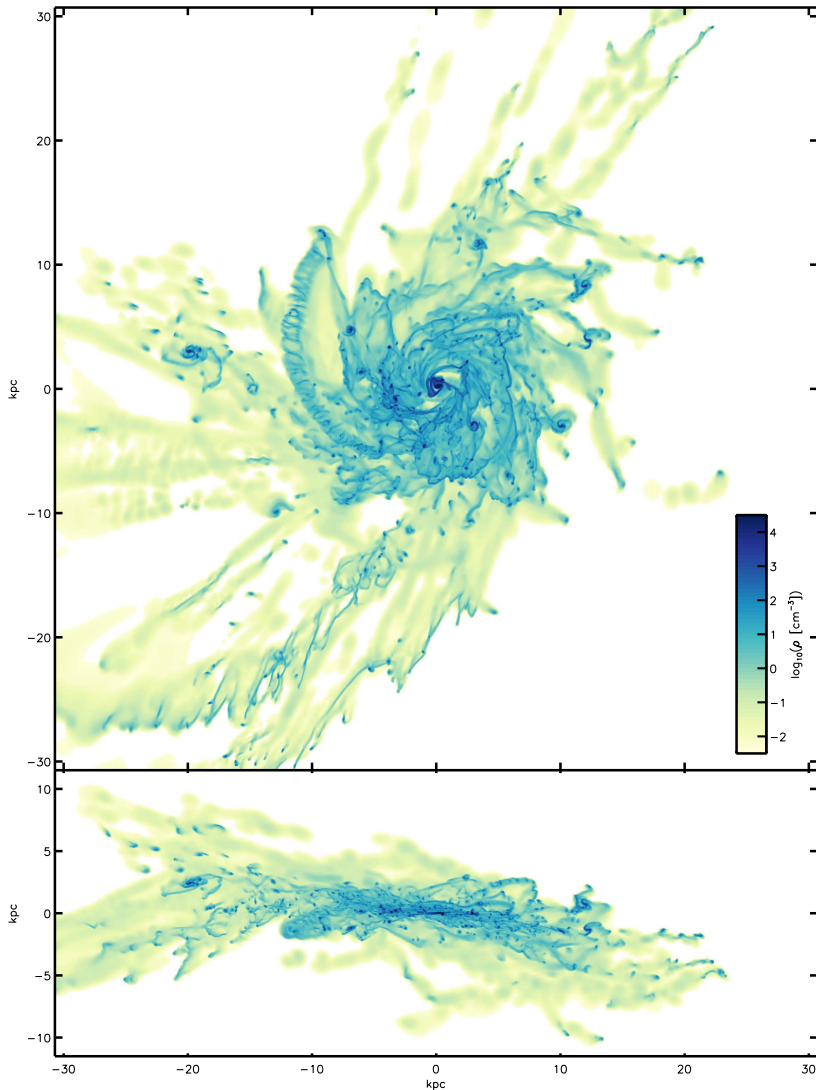


Fig. 4. Face-on (top) and edge-on (bottom) mass-weighted average density maps of the gas for the G1_G1_90_90_0 simulation 280 Myr after the coalescence (i.e., 640 Myr of evolution after the initial conditions).

the MIRAGE sample, covering the evolution up to 800 Myr, displayed in 11 time steps. The whole MIRAGE sample maps, containing 23 figures, are available as online material.

5.1. Initial conditions relaxation

The relaxation of the disk plays a fundamental role at the beginning of the simulation. The low halo concentration when compared to lower redshift, combined with a high gas fraction drives the gas disk towards an unstable state with $Q < 1$, even though we start our simulation with the requirement $Q > 1.5$ everywhere in the stellar disks. The high cooling rates of the gas in the initial disk allow very fast dissipation of internal energy. To prevent uncontrollable relaxation that occurs rapidly, we start our simulations with a maximum resolution of 59 pc ($\ell_{max} = 12$) and with a temperature floor for the gas of $T = 10^4$ K (see Table 4). To establish the turbulence smoothly afterwards, we pro-

gressively increase the resolution every 25 Myr starting from 85 Myr, until we reach a maximum resolution of 7.3 pc, and a temperature floor for the gas of 300 K. This allows the disks spiral features supported by the thermal floor to form quickly during the first time steps. Once the resolution is increased, Jeans instabilities arise and give birth to clumps of a few 10^8 solar masses, which can quickly merge to form more massive ones. We observe a rapid contraction of the disks, reducing their radial size by $\sim 20\%$ during the first 80 Myr (about a third of the dynamical time), owing to the dissipation of energy by the gas component. This ad-hoc relaxation strategy insures that the internal energy of the gas disk dissipates gradually through cooling over 130 Myr, and also helps us to save computational resources. The refinement down to the level $\ell = 14$ at $t = 105$ Myr allows reaching densities $\rho > \rho_0$, thus enabling the formation of stars and all the associated feedback of newly formed stars.

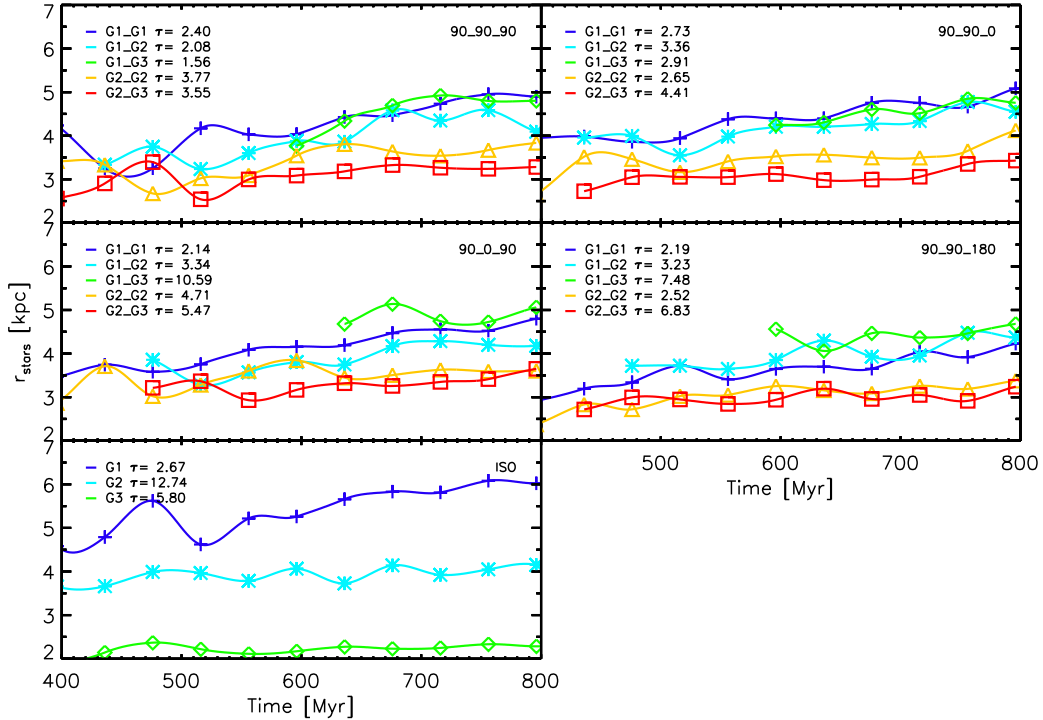


Fig. 5. Evolution of the stars disk scalelength in the MIRAGE sample. Each panel traces the evolution of the scalelength for a given orbital configuration, allowing a comparison between mass ratios for a given set of disk orientations at a given specific orbital energy. The measurements are performed every 40 Myr, starting at the time of the core coalescence (400 Myr for the fastest mergers), and each curve linking these measurements is the result of a cubic interpolation to increase the clarity of the plot. The colored lines and different symbols indicate the mass ratio of the progenitors (given by $G_i G_j$, see Table 3); the label at the top right of each panel indicates the initial orientation of the disks (given by $\theta_1 \theta_2 \kappa$, see Table 2). The lower left panel is dedicated to isolated simulations. For each simulation, we indicate the growth time τ expressed in Gyr, which is the time needed for the disk/remnant to double its size starting from the closest measurement to 400 Myr.

t [Myr]	ℓ_{max}	Δx [pc]	T_{floor} [K]
[0, 80]	12	58.6	10^4
[80, 105]	13	29.3	10^3
[105, 130]	14	14.6	500
[130, 800]	15	7.3	300

Table 4. Refinement strategy of the high redshift disks.

5.2. Gas accretion from hot halo

As mentioned in section 4.3, the AMR box is continuously replenished with metal-free gas. The very low-density component is constrained by a gamma-polytrope, which ensures the formation of a hot stabilized halo. The central part of the halo reaches densities above 10^{-3} cm^{-3} , where the pressure support from the gamma-polytrope ends. Thanks to metal lines cooling, the central part of the gaseous halo can cool down and condense on the galactic disk. We measure the accretion rates in a spherical shell of 20 kpc (typical value of the halo scalelength in the most massive galaxy model), by detecting the metal poor gas ($Z < 10^{-3}$) able to enter the sphere within a time step of 5 Myr. In Table 5 we display the mean values of these accretion rates, as well as the mean SFR for the different masses configurations of the sample. We compare these values to theoretical predictions from the baryonic growth rate formula found in Dekel et al. (2009a). The theoretical values are obtained using total halo masses (without

considering radial cuts), and we assume that only two thirds of this accretion rate can be associated to smooth gas accretion, the remaining third being associated to mergers as observed in Dekel et al. (2009a).

	$\langle \dot{M}_{gas} \rangle$ [$M_{\odot} \text{ yr}^{-1}$]	$\langle \text{SFR} \rangle$ [$M_{\odot} \text{ yr}^{-1}$]	$\dot{M}_{th}(z=2)$ [$M_{\odot} \text{ yr}^{-1}$]	$\dot{M}_{th}(z=1.5)$ [$M_{\odot} \text{ yr}^{-1}$]
G1	13.4	17.5	31.8	21.1
G2	2.6	7.7	11.0	7.3
G3	0.8	3.7	3.8	2.5
G1_G1	41.4	31.5	63.6	42.2
G1_G2	20.8	23.2	42.8	28.4
G1_G3	17.4	20.2	35.6	23.6
G2_G2	8.0	13.9	22.0	14.6
G2_G3	6.0	10.7	14.8	9.9

Table 5. Comparison of the mean star formation and accretion rates measured in the MIRAGE sample. $\langle \dot{M}_{gas} \rangle$: average accretion rate of inflowing pristine gas ($Z < 10^{-3}$) for the isolated disks and the mergers measured in a spherical shell with a radius of 20 kpc. $\langle \text{SFR} \rangle$: average SFR. $\dot{M}_{th}(z)$: Theoretical prediction of the gas accretion rate as function of redshift and halo mass. All the averages are computed in the interval [100,800] Myr.

Furthermore, at $z=2$, Agertz et al. (2009) find an accretion rate of hot gas for a galaxy with a baryonic mass $M_b \sim 10^{11} M_{\odot}$

in good agreement with our model having a close mass (namely the G1 model). They also show that the cold gas accretion that prevails at $z \gtrsim 2$ becomes dominated by hot gas accretion at lower redshifts, which makes our implementation agree with this statement. This scenario is also supported by recent work that uses moving-mesh codes, which find a substantially lower cold gas accretion rate than in comparable SPH simulations (Nelson et al. 2013). The gas accretion rate in the MIRAGE sample slightly increases with time (see section 4.3), implying that simulating more than 1 Gyr of evolution would lead to unrealistic high accretion rates. The mean accretion rates measured over 800 Myr in our simulations remain consistent with theory and cosmological simulations (see Table 5).

5.3. Mass-size evolution

To follow the evolution of the mass-size relation in our simulation sample, we proceed to a centering and spin alignment with the z -axis of the AMR cartesian grid. As the stars age, they experience a progressive gravitational heating that redistributes the oldest stars into a diffuse halo component with a smoother gravitational potential. The center of the each simulation is therefore found by using the peak of the mass-weighted histogram of the positions of the oldest stars, i.e. those stars only present in the initial conditions. This peak in the distribution of the old stars is hereafter associated to the center of the bulge. We recover the galactic disk orientation using the spin vector of the stars younger than 50 Myr for a given snapshot. They tend to still be located within the gas disk, which cannot be used to perform such a computation because of the turbulence and the outflows carrying consequent momentum. Once the orientation of the galactic disk is correctly recovered, we compute the stellar surface density profile. The distribution is decomposed into a bulge and a disk by performing a linear regression on the surface density profile in the interval $[r_{cut,bulge}, r_{cut,stars}]$ to extract the disk profile (see Table 1). To each surface density measurement we associate a relative error proportional to the square root of the number of particles found within the radial bin.

The evolution of the disk scalelengths is displayed in Fig. 5. It confirms that both mergers and isolated disks can produce an inside-out growth (Naab & Ostriker 2006) regardless of the orbital configurations, and despite the proven ability of gas-rich mergers to produce compact systems (Bournaud et al. 2011). For each simulation, we estimate the growth time, which we define as the time needed for the stellar disk to double its size measured right after the coalescence (or at $t=400$ Myr for the isolated systems). A mean growth time of 3.9 Gyr is measured for the mergers, with the fastest systems reaching growth time close to 2 Gyr. It appears that the less massive systems, therefore the less clumpy, are less efficient at driving stellar disk growth. This inside-out growth is taking place in an idealized framework, although the galaxies are accreting gas from the halo at a rate comparable to cosmological simulations (see Section 5.2). This continuous gas accretion fuels secular evolution processes that are able to drive such growth by performing a mass redistribution. Consequently, our results suggest that other mechanisms than late infall of cold gas from the cosmic web (Pichon et al. 2011) may alternatively build up high redshift disks inside-out.

The stellar mass-size relation for the MIRAGE sample is shown in Fig. 6. We plot the mass-size relation found in Dutton et al. (2011) and shifted at different redshifts. Our choice of stellar sizes in the initial conditions makes the simulations of the MIRAGE sample lie in the dispersion range computed for the $z = 1.5$ mass-size relation. Nonetheless, the size evo-

lution is fast in the MIRAGE sample, although one can expect this rapid growth to stop once the clumpy phase ends. Indeed, the size growth is linked to the gas-rich clump interaction that is able to redistribute significant amount of stellar mass toward the outskirts of the disk. We overplot on the simulations data the values for the MASSIV sample, and the error bars show the 1σ standard deviation computed using the errors on the stellar mass and size (Vergani et al. 2012). We use the classification of López-Sanjuan et al. (2013) to differentiate isolated galaxies from minor and major mergers on the plot using different symbols. We observe that the majority of galaxies classified as major merger lie above or close to the $z = 1.5$ mass-size relation, which is straightforward once one considers that the size measurement is done on an extended system where the two disks are not yet mixed well. This gives credit for the major merger classification performed by López-Sanjuan et al. (2013). Overall, the bulk of the MASSIV sample ranges within the dispersion fork of the $z = 1.5$ relation, which makes our simulations consistent with observations. A fraction of the isolated and minor merger systems are more than 1σ below the $z = 1.5$ relation, suggesting a population of compact galaxies.

5.4. Star formation

Figure 7 presents the star formation histories of the MIRAGE sample for the different masses and merger orbital configurations. For each simulation, we indicate the coalescence time t_c expressed in Myr in the legend, as well as the pericentral time t_{peri} . The star formation histories exhibit stochastic behavior due to the clump interactions and the cycling energy injection by stellar feedback maintaining the gas turbulence. The mean ratio of the SFR dispersion over the average SFR ($\sigma_{SFR}/\langle SFR \rangle$) for the whole MIRAGE sample is roughly equal to 30%. Quite surprisingly at first glance, we do not observe any SFR enhancement due to the galaxy merger. Neither orientations nor mass configurations appear to produce enhanced SFR.

Figure 8 shows the histogram of the normalized quantity in the interval $t_c \pm 100$ Myr:

$$(\text{SFR}(t) - \text{SFR}_{\text{iso}}(t)) / \langle \text{SFR}_{\text{iso}} \rangle$$

with SFR_{iso} the summed SFR of the fiducial simulations evolved in isolation, and $\langle \text{SFR}_{\text{iso}} \rangle$ the mean value of SFR_{iso} . For each histogram, we display the value of barycenter of the distribution β in the legend, which allows estimating how much the interaction enhances the star formation in the time interval defined previously. We observe no trend to SFR enhancement due to merger ($\beta \leq 0$), even if in the case G1_G3 two of the merger produces somewhat more stars than the summed fiducial isolated models ($0.1 \leq \beta \leq 0.2$). However, this value is too low to be considered as a starburst. Generally, the mergers are even less effective at producing stars compared to isolated simulations. This result contradicts other works (e.g. Bournaud et al. 2011; Teyssier et al. 2010; Cox et al. 2008; Powell et al. 2013). Since this paper does not intend to perform a full study of the starburst efficiency in high redshift galaxies, we list subsequently and briefly discuss the possible reasons for the suppression of starburst in our simulation sample:

- Our choice of elliptic Keplerian trajectories might affect the star formation efficiencies of our merger simulations. However, many works have demonstrated that the starburst efficiency of equal mass galaxy mergers is insensitive to the orbits, the disk orientations, and the physical properties of these galaxies (e.g., Mihos & Hernquist 1996; Springel 2000;

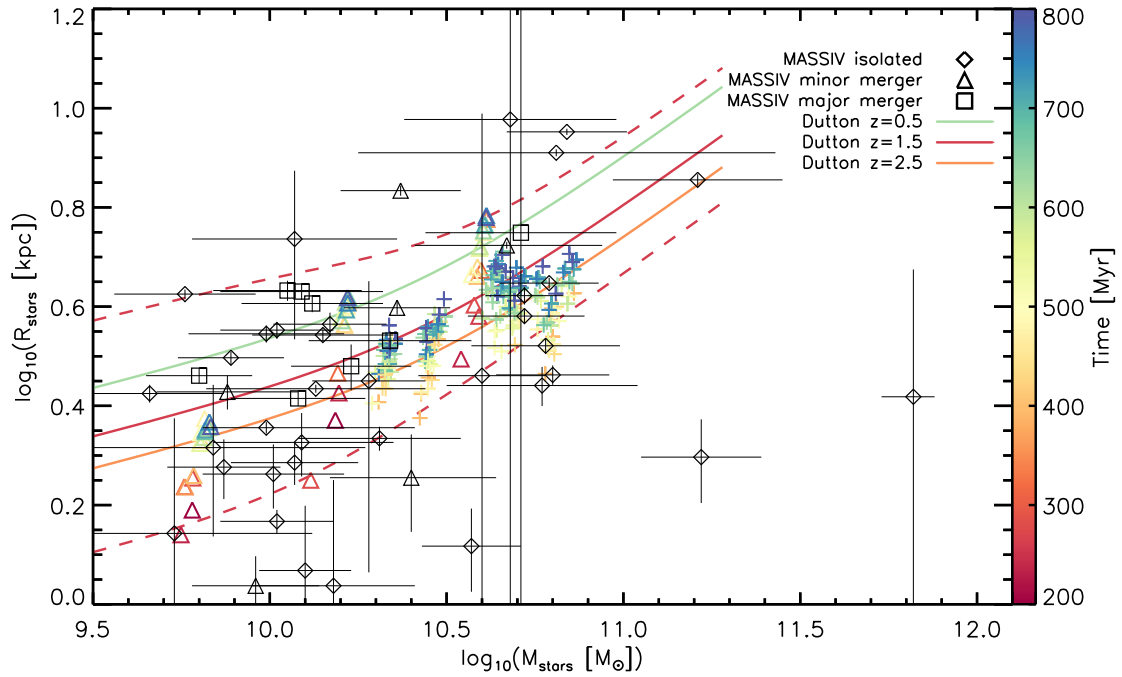


Fig. 6. Stellar mass as a function of stellar scalelength. The symbols “+” and “ Δ ” show the MIRAGE galaxy mergers and isolated disks, respectively. The color encodes the time evolution since the initial conditions. Black symbols display the MASSIV data, according to measurements found in Vergani et al. (2012) and Epinat et al. (2012). The stellar mass-size relation derived in Dutton et al. (2011) and shifted to $z = 1.5$ is overlotted with the red solid curve. The dotted curves show the dispersion computed for $z = 1.5$ from the relation derived in Dutton et al. (2011). We also display the mass-size relation for $z = 0.5$ (green line) and $z = 2.5$ (orange line) to emphasize the redshift evolution of the relation.

Cox et al. 2004). Consequently, the initial configuration of the major merger simulations G1_G1 and G2_G2 should not be considered as responsible for the absence of starburst. However, longer interactions would lead to coalescence of more concentrated systems because of the clumps migration, which could enhance a nuclear starburst. That higher mass ratios simulations (G1_G2, G1_G3, G2_G3) that explore more elongated orbits (see Table 3) do not exhibit star formation enhancement might suggest that the orbits and the disk orientations are generally not to blame for this lack of star formation overactivity.

- As highlighted by Moster et al. (2011), the hot gaseous halos implied in a galaxy merger are likely to be heated by shocks, together with an acquisition of specific angular momentum increasing the centrifugal barrier. Both of these processes can push toward a lower starburst efficiency because isolated disks are more effective at accreting gas from the hot halo.
- A complex treatment of the ISM favors the production of hot gas, which systematically lowers SF, as Cox et al. (2006) point out. The simulations performed in Bournaud et al. (2011) constitute a good dataset for direct comparison, owing to the initial conditions definitions very close to our G1 model. The presence of starbursts in comparable simulations when the gas obeys to a 1D equation-of-state suggests a change in the gas response to a galactic interaction.
- Teyssier et al. (2010) demonstrate that the starburst in a low redshift major merger is mostly driven by the enhancement of gas turbulence and fragmentation as long as the numerical resolution allow it to be resolved. It may be more difficult to increase this turbulence and fragmentation at high redshift

because both are already high in our isolated gas-rich disks. The isolated disks simulations are indeed able to maintain this high level of turbulence and fragmentation thanks to continuous gas refilling by the hot halo accretion and an efficient stellar feedback. This scenario would suggest that star formation can saturate and prevent starbursts in galaxy mergers of very turbulent and clumpy gas-rich disks.

- High gas fractions ($>50\%$) are maintained throughout the duration of the mergers. These high gas fractions may prevent the formation of a stellar bar in the remnant, which would drag a large amount of gas toward the nucleus to fuel a starburst (Hopkins et al. 2009). The large fraction of cold fragmented gas prevents the formation of a bar in the stellar component. Additionally, the stellar feedback removes gas from the star-forming regions continuously and may also act against the formation of a large stellar disk by lowering the SFR Moster et al. (2011).
- The feedback model adopted in this study might not be energetic enough to succeed in ejecting important quantities of gas on very large scales, especially because of the isotropic hot gas accretion that systematically curbs the outflowing material. The adopted feedback model may be efficient enough to saturate the star formation during the pre-merger regime, but is not strong enough to deplete the disk from large quantities of gas, which would then be accreted again later, feeding a star formation burst.

Numerous processes can explain the starburst removal in very gas-rich clumpy and turbulent galaxy mergers. The star formation histories of the MIRAGE sample remain difficult to interpret without a complete study in a full cosmological environment

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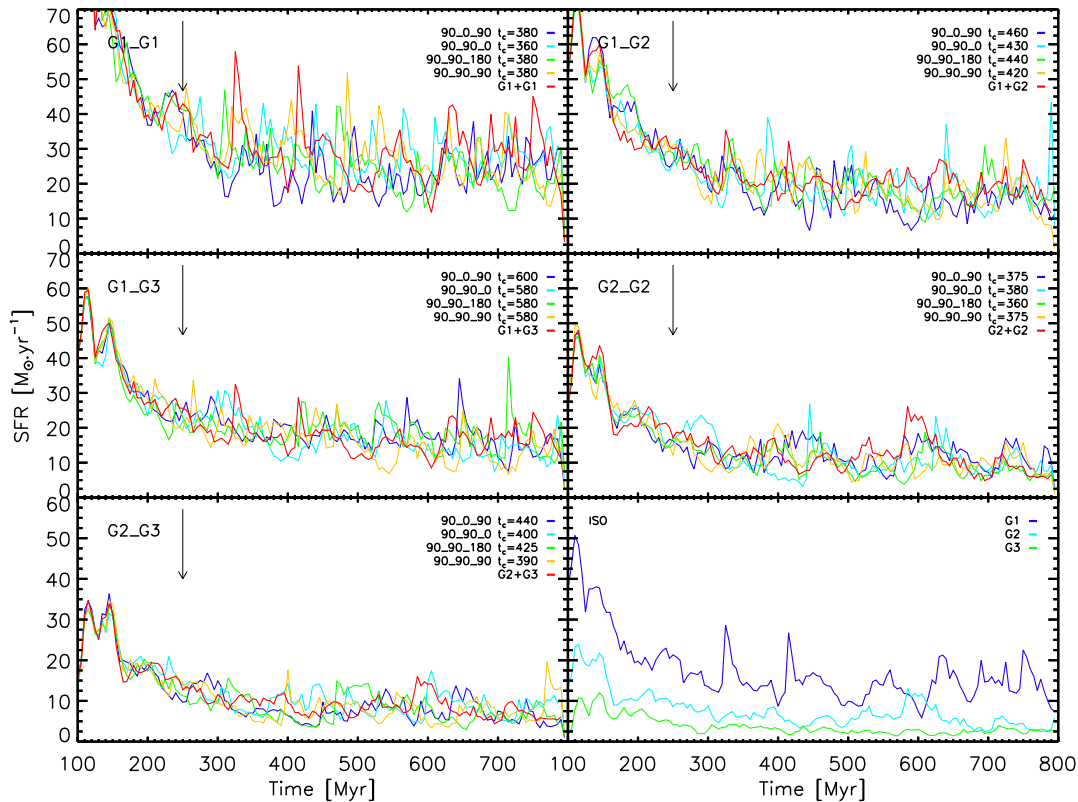


Fig. 7. Star formation histories for each simulation of the MIRAGE sample. Each panel explores disk orientations for fixed masses respectively given by $\theta_1\theta_2\kappa$ and G_iG_j (written on the top right of each panel, see Tables 2 and 3). The last panel shows the SFR of the isolated disk simulations. The curves begin at 100 Myr (see section 5.1). To compare the SFR of merging disks with the SFR of isolated disk per mass unit, the SFR of isolated disks (red dotted lines) have been superimposed on the SFR of merging disks. The black arrow in the merger panels shows the pericentral time t_{peri} equal to 250 Myr in all the merger simulations. For each galaxy merger, we also display the time of the coalescence of the galactic cores t_c visually determined.

to weight each configuration according to its occurrence probability. Generally, the link between mergers and starburst may be more fuzzy at high redshift than at lower redshift.

Figure 9 displays the SFR as a function of the stellar mass. We compared the MASSIV “first-epoch” data with the MIRAGE sample, for which the SFR has been estimated from the integrated $H\alpha$ luminosity, and stellar mass within the optical radius. The MASSIV error bars were computed using the errors on the $H\alpha$ flux measurement found in Queyrel et al. (2012). The scatter observed for a given simulation stellar and gas mass underlines the stochastic nature of the star formation in gas-rich clumpy disks. This scatter is nevertheless still lower than the one observed in the MASSIV data, which encompasses very varied gas fractions.

Figure 10 displays the position of the MIRAGE sample on the KS diagram between 200 Myr and 800 Myr for the isolated disks and between the coalescence and 800 Myr for the merger simulations. We computed the gas surface density Σ_{gas} and the star formation surface density Σ_{SFR} quantities within r_{stars} , the stellar disk scalelength estimated with the method described in section 5.3. We also rejected all the gas in the IGM by only considering cells with densities greater than $\rho = 2 \times 10^{-3} \text{ cm}^{-3}$, which typically corresponds to the frontier between ISM and

IGM in all of the simulations. The quantities Σ_{gas} and Σ_{SFR} are measured on face-on projections, after having centered our referential on the peak of the old stars probability distribution function and aligned the spin of the young stars disks with our LOS. We note that our sample lies close to the relation found in Daddi et al. (2010b), with a slight shift towards lower star formation efficiencies (within the 1σ dispersion), which can be attributed to the shutdown of star formation at high gas temperatures. The MIRAGE sample does not show any bimodality, as expected from the star formation histories displayed in Fig. 7. One should take into account that, by construction, our sample does not provide the statistical cosmological weight of a volume-limited sample of the $1 < z < 2$ galaxy population.

We also overplot the position of the MASSIV sample on this diagram for comparative purposes. The MASSIV error bars on the quantity Σ_{SFR} are computed again using the uncertainties on the $H\alpha$ flux. We also take an error proportional to the spatial sampling of the SINFONI data into account, which we propagate to the measurement of the radius of the ionized emission region. We do not have a strong observational constraints for the amount of gas in the MASSIV galaxies. Nevertheless, we mark out the gas mass for each galaxy assuming a mean gas fraction $f_g = 45\%$, which is the mean value obtained on the dynamical

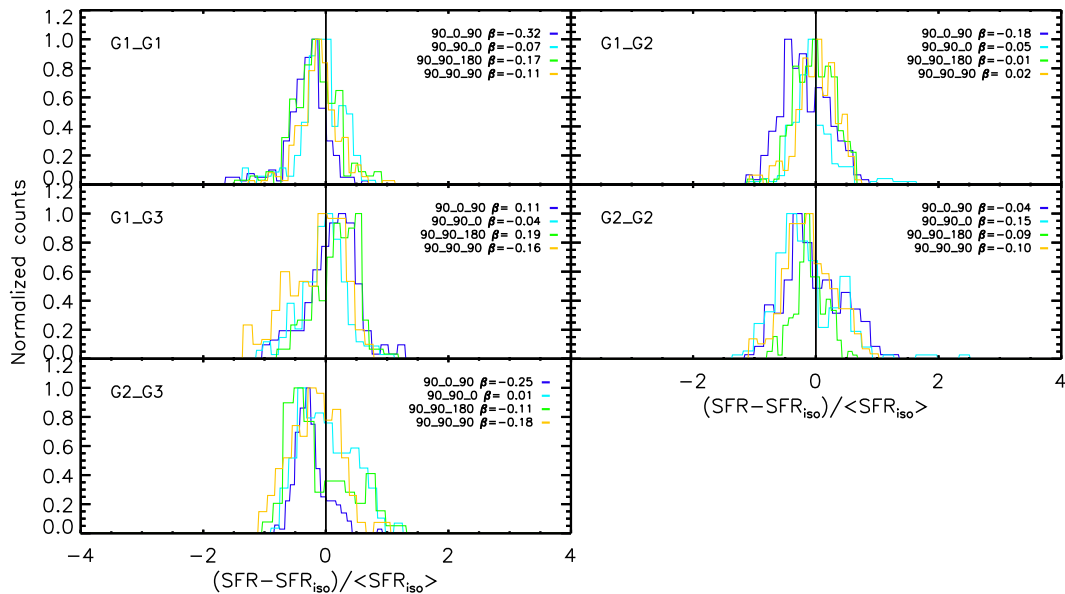


Fig. 8. Histogram of the difference between the merger SFR and the cumulative isolated SFR ($SFR - SFR_{iso}$) computed between 100 Myr before and 100 Myr after the galaxies coalescence. Each panel explores disk orientations for fixed masses respectively given by $\theta_1, \theta_2, \kappa$ and G_i, G_j (written on the top left of each panel, see Tables 2 and 3). From these histograms, we can interpret how much time a merger spends with a higher or lower SFR during this crucial period. The quantity β in the legend is the barycenter of the histogram, which measures the shift in star formation induced by the merger compared to secular evolution.

cal/stellar mass diagram of the MASSIV sample (Vergani et al. 2012). Using the relation $M_{gas} = f_g M_{stars} / (1 - f_g)$, we can overplot the MASSIV data on the KS diagram. We compute the errors bars of the Σ_{gas} quantity assuming a minimal gas fraction of $f_{g,min} = 25\%$ and $f_{g,max} = 65\%$, a range where we can expect the MASSIV sample to lie. We then propagate the errors on the stellar mass using $f_{g,min}$ and $f_{g,max}$. Therefore, the distribution of the MASSIV data on the KS relation is close to the “normal” regime of star formation, considering our assumptions on the gas fractions. Our merger simulations match the area covered by both the isolated and merging galaxies of the MASSIV sample on the KS diagram.

6. Summary and prospects

In this paper, we introduce a new sample of idealized AMR simulations of high redshift ($1 < z < 2$) mergers and isolated disks referred to as MIRAGE. The sample is originally designed to study the impact of galaxy merger on the gas kinematics in a clumpy turbulent medium. We focus on presenting of the methods used to build the MIRAGE sample and on the first results obtained for the evolution of the masses, sizes, and star formation rates.

The key points of the goals and methods used in this paper can be summarized as follows:

- We presented the MIRAGE sample, a series of mergers and isolated simulations using the AMR technique in an idealized framework that compare extreme signatures in terms of gas kinematics. The MIRAGE sample initial conditions probe four disk orientations (with κ ranging from 0° to 180°), five total baryonic merger masses (ranging from 4.9 to $17.5 \times 10^{10} M_\odot$) and three galaxy mass ratios ($1:1, 1:2.5, 1:6.3$)

among 20 merger simulations designed from three disk models (with baryonic masses of $1.4, 3.5,$ and $8.8 \times 10^{10} M_\odot$). The case of low gas fractions has been extensively studied in the literature, so we choose here to only study gas-rich galaxies ($f_g \sim 60\%$) to study the impact of the presence of giant star-forming clumps on merging turbulent disks.

- We introduced DICE, a new public code designed to build idealized initial conditions. The initialization method is similar to what has been done in Springel et al. (2005b). The use of MCMC algorithm to build a statistical distribution requiring only the 3D-density function as input allows us to consider building components in future developments with more complex density functions compared to the canonical ones used in this paper.
- We used a new implementation of stellar feedback from the young, massive part of the IMF (Renaud et al. 2013), coupled to a supernova feedback with non-thermal processes modeled using a cooling switch (Teyssier et al. 2013). The new physically-motivated implementation of young-star feedback allowed us to track the formation of Strömgren spheres where the energy from the massive young stars is deposited, allowing future comparisons with simulations using feedback recipes parametrized with wind mass-loading factors.

The key results of this paper can be summarized as follows:

- Star formation in disks – We find that the star formation history of isolated disk galaxies fluctuates strongly throughout the duration of simulations, with a SFR dispersion close to 30% around its mean value. This star formation proceeds mostly in giant clumps of gas and stars and naturally gets a stochastic behavior. The small star formation bursts may account for the intrinsic scatter of the “main sequence” of star forming galaxies at $z = 1 - 2$ (Daddi et al. 2010b).

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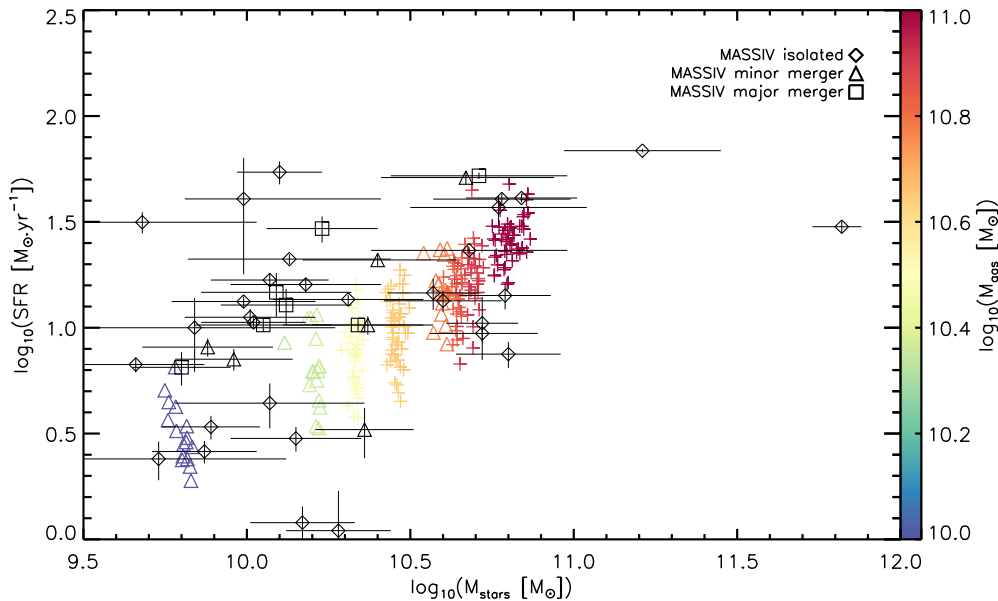


Fig. 9. Star formation rate as a function of stellar mass measured between the coalescence and 800 Myr for the merger simulations, and between 200 and 800 Myr for the isolated simulations. Black symbols show MASSIV data for which the SFR is estimated from the H α integrated luminosity, and the stellar masses measured within the optical radius $r_{opt} = 3.2 \times r_{stars}$. Each colored symbol shows a snapshot of the MIRAGE mergers and isolated disks simulations, respectively plotted with “+” and “ Δ ”. The color encodes the gas mass of the disks and remnants measured within the gas optical radius.

- Star formation in mergers – The minor and major gas-rich mergers of our sample do not induce major bursts of star formation significantly greater than the intrinsic fluctuations of the star formation activity. The mechanisms for triggering active starburst at high redshift could be different from the ones at low redshift due to large differences in the amount of gas available for accretion in the circumgalactic medium around stellar disks. This suggests that a complex modeling of the gas capturing a high level of fragmentation and turbulence maintained by stellar feedback and gas accretion may offer a mechanism of saturation for the star formation activity in high- z galaxies. The remarkable homogeneity of the observed specific SFR in high redshift galaxies (Elbaz et al. 2007, 2011; Nordon et al. 2012), coupled to the prediction of a high occurrence of minor mergers in this redshift and mass range (Dekel et al. 2009a; Brooks et al. 2009), may support the scarcity of star formation bursts triggered by very gas-rich mergers.
- Star formation scaling laws – Overall, our sample of disks and mergers is compatible with the evolution of the mass-SFR relation observed for a complete sample of star-forming galaxies in the same mass and redshift range (namely the sample MASSIV, Contini et al. 2012), and independently of the assumed disk and merger fraction in the sample. In a Kennicutt-Schmidt diagnostic, the majority of mergers are close to the “normal” regime of disk-like star formation as defined by Daddi et al. (2010b), Genzel et al. (2010), with a slight deviation towards lower star formation efficiencies.
- Size evolution – A stellar mass-size relation in accordance with Dutton et al. (2011) has been obtained by our models, and the evolution with redshift of this relation was also reproduced. In particular, inside-out growth can be obtained as a natural outcome of internal dynamical processes redistributing angular momentum mostly through clumps interactions:

these processes can naturally make disks become larger over time, for any given stellar mass, even if mergers are expected to produce more compact systems. Our simulations only include infall of low-angular momentum material through hot gas accretion, suggesting that the radial inside-out growth at the observed rate might not need to be achieved through a cold mode in the context of our idealized modeling.

The significant fraction of active galactic nucleus (AGN) in the redshift range $1 < z < 2$ means that the inclusion of black-hole particles and the associated AGN feedback should be addressed specifically. Nevertheless, the recent work of Gabor & Bournaud (2013) shows that AGN feedback is unable to disrupt the clumps of 10^8 - $10^9 M_{\odot}$ formed *in-situ* in comparable idealized gas-rich disk simulations. Dubois et al. (2013) also show that massive clumps may survive AGN feedback during their migration towards the bulge in a fully cosmological context. Since the clumps drive most of the SF, we do not expect major changes on short term star formation histories by including AGN feedback. However, at later stages of evolution, the strong heating of the gaseous halo driven by shocks due to AGN feedback should lower the accretion rate and lead to lower gas fractions in the merger remnants. Finally, the results obtained in the MIRAGE sample call for further investigations to assess the effect of AGN feedback in this kind of simulations.

With the MIRAGE sample we have the opportunity to investigate further questions of galaxy evolution. The combination of statistical probing of the orbital parameters, the controlled input parameters due to idealized framework, the parsec scale resolution, and the explicit physically motivated implementation of stellar feedback make it a useful database to use for studying (i) the properties and lifetime of the giant (10^8 - $10^9 M_{\odot}$) star-forming clumps (already presented in Bournaud et al. 2013), (ii) the impact of the migration and interaction of the clumps on the

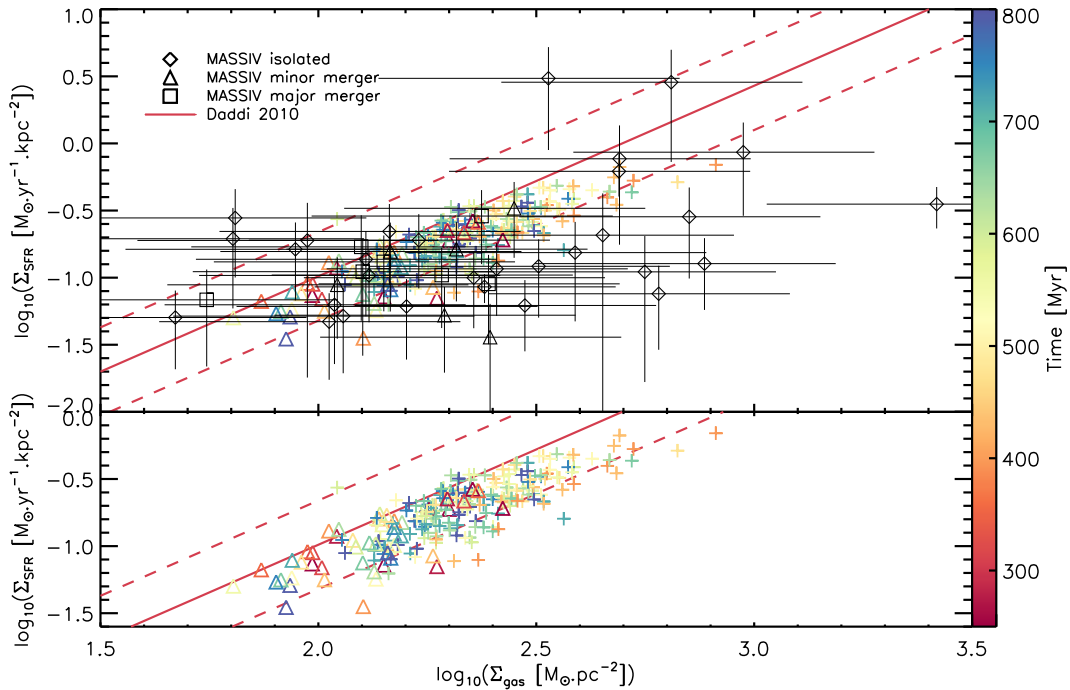


Fig. 10. Kennicutt-Schmidt relation for the simulations involved in this study. We use two panels for clarity; in the top panel we overplot the MASSIV data on the MIRAGE sample for comparison. In both panels, we also display the relation obtained in Daddi et al. (2010b) (red solid line for the relation and dashed line for the associated 1σ dispersion). Simulations plotted at different times are represented with different colors, with values measured inside the stellar disk scalelength. In case of merger, we plot only snapshots where the coalescence has been reached. The mergers and isolated disks are respectively plotted with “+” and “ Δ ”. The MASSIV sample (Contini et al. 2012) positions are computed using the half-mass stellar radius for a typical gas fraction of 45%, and are plotted using black diamonds for isolated galaxies, triangles for minor mergers, and squares for major mergers. The associated error bars are computed using the errors on $H\alpha$ flux, stellar size, and the assumption that the gas fraction f_g lies in the range $[0.25, 0.65]$.

galaxies properties, (iii) the metallicity evolution in mergers and isolated disks, (iv) the classification of velocity fields of high- z galaxies based on a large set of mock observations derived from the MIRAGE sample, among other studies.

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Granularité des galaxies juvéniles 4

Résumé

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L'analyse morphologique des galaxies à grand décalage spectral révèle une fraction importante de systèmes où la matière est distribuée de manière asymétrique, sous forme de granules massifs composés de gaz et d'étoiles. Dans ce chapitre, je passe en revue les différentes observations de galaxies à grand décalage spectral, et les principales conclusions associées. Après avoir décrit le mécanisme de formation de ces granules dans les disques jeunes et riches en gaz, je présente ma contribution aux questions relatives à l'évolution de ces granules grâce à l'analyse de l'échantillon des simulations MIRAGE. Un second article concentré sur l'analyse des granules dans des simulations de disques isolés est présenté. Enfin, la dernière partie de ce chapitre présente une analyse globale du mécanisme de coalescence de granules et ses effets sur l'évolution à long terme sur les disques isolés et les rémanents de fusion.

4.1 Observations

Dans les galaxies spirales de l'Univers local, les régions de formation stellaire ont des masses qui excèdent rarement 0.1% de la masse totale du disque. L'ensemble des régions de formation stellaire d'une galaxie est donc loin de dominer la distribution totale de masse. L'avènement des observations de champs profonds à grande résolution spatiale a permis de révéler que la majorité des galaxies "normales" formant des étoiles à grand décalage spectral ($z > 1$) avaient une morphologie asymétrique où la lumière est confinée dans des granules¹ possédant des tailles de l'ordre du kilo-parsec, et des masses comprises entre 10^8 et $10^9 M_{\odot}$ (Fig 4.1 et 4.2). Ces galaxies ne sont caractérisées que rarement par des structures spirales. Contrairement aux galaxies locales, ces granules peuvent contenir une grande fraction ($\sim 40\%$) de la lumière totale émise par les galaxies (Elmegreen et al., 2005). Historiquement, on distingue deux types de morphologies granulaires : les "galaxies en chaîne"² (Cowie et al., 1996) qui exhibent une série de granules alignés sur un même axe (voir les vignettes C1, C8, C9 de la figure 4.2), et les "amas granulaires"³ ressemblant à des disques dont la distribution lumineuse est concentrée dans quelques granules (voir toutes les autres vignettes de la figure 4.2). Dans les deux classes, les granules ont des propriétés physiques similaires, suggérant que les galaxies granulaires en chaîne seraient identiques aux amas de granules, mais simplement observées à des inclinaisons différentes. La structure de ces galaxies rappelle celle des galaxies subissant une fusion dans l'Univers proche, d'où une certaine incertitude quant à la véritable nature de ces objets. Cependant, on ne distingue pas dans toutes les galaxies granulaires des structures typiques des interactions telles que les bras, les débris de marées, ainsi que les anneaux, suggérant que la plupart de ces objets sont en fait des galaxies isolées. Les galaxies granulaires de l'Univers lointain ressemblent aux galaxies naines irrégulières de l'Univers local en termes de morphologies (figure 4.3), néanmoins les galaxies granulaires possèdent des luminosités et des masses similaires à celles des galaxies spirales, ce qui exclut tout lien de parenté entre naines irrégulières et galaxies granulaires. Elmegreen et al. (2009a) notent que les bulbes présents dans les galaxies granulaires sont beaucoup moins massifs et beaucoup plus jeunes que leurs analogues locaux, ce qui renforce l'idée que les galaxies granulaires sont des galaxies spirales jeunes dans un état de transition instable où le bulbe stellaire commence seulement à se former.

Elmegreen et al. (2009b) notent la présence d'une évolution depuis les galaxies granulaires sans disque sous-jacent vers des amas granulaires ayant une composante stellaire rouge sous forme de disque, avec au centre un bulbe qui se distingue de manière claire des granules. Ces observations soulignent que cette évolution est donc probablement le fruit de processus internes, durant lesquels les granules interagissent entre eux tout en migrant progressivement vers le centre de masse, amenant ainsi de la masse permettant de construire petit à petit le bulbe stellaire. Les disques à grand décalage spectral formeraient donc des étoiles dans des complexes très massifs résultant d'instabilités gravitationnelles majeures dans leur composante gazeuse. Ces granules formés par instabilités sont trois fois plus grands et ~ 100 fois plus massifs que les complexes de formation stellaire des galaxies locales. Cependant, ces deux types de complexes semblent être tous les deux définis par la longueur et la masse de Jeans. Cette augmentation de la taille des complexes de formation stellaire peut engendrer un facteur ~ 5 dans les vitesses turbulentes et un facteur ~ 10

1. En anglais, ces granules sont désignés par le terme "clump". De plus, on désigne une galaxie granulaire par le terme "clumpy galaxy".

2. "Chain galaxies" en anglais.

3. "Clump clusters" en anglais.

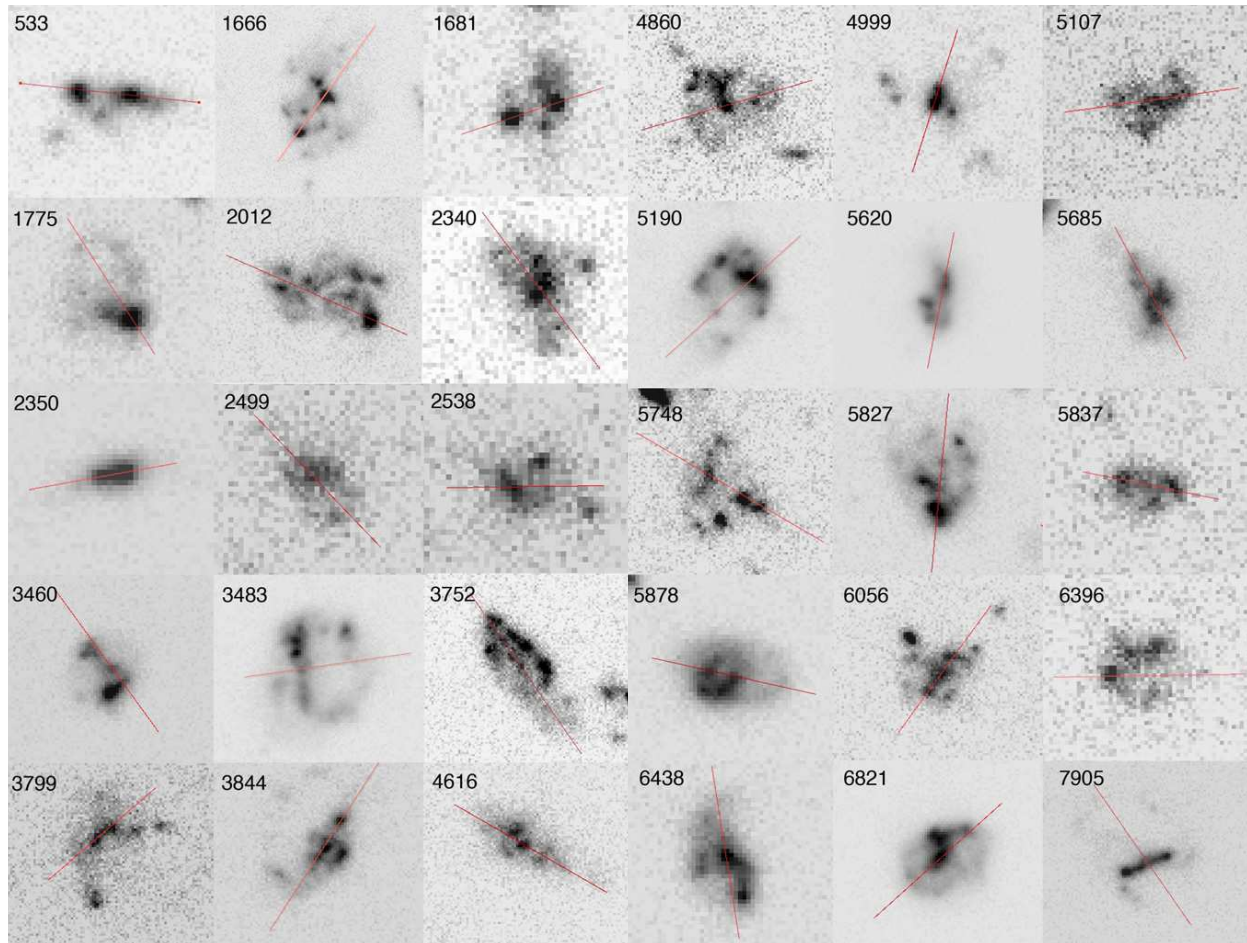


Figure 4.1 – Images des galaxies granulaires détectées dans le Hubble *Ultra Deep Field*, observées dans la bande passante $i775$ avec la caméra ACS du télescope Hubble. L'échantillon présenté dans cette figure possède un décalage spectral moyen de $z=1.7$. Crédits : Elmegreen et al. (2013).

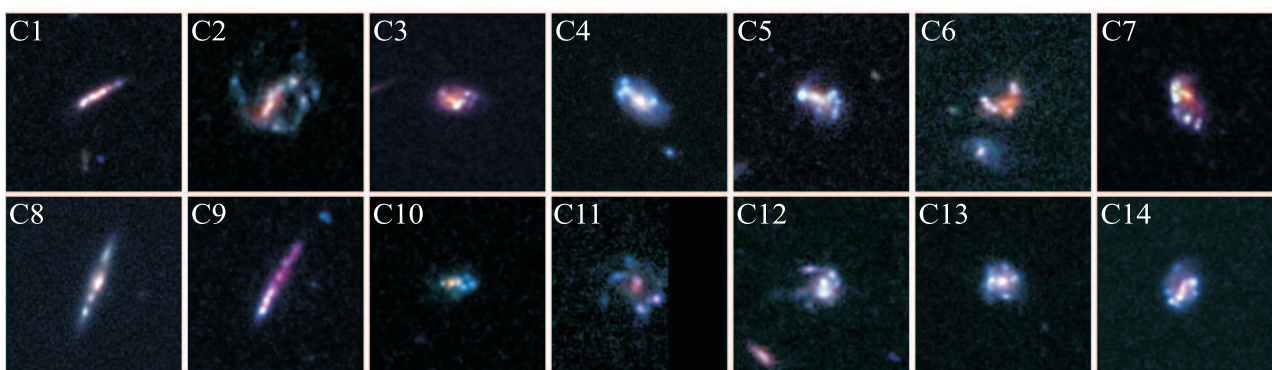


Figure 4.2 – Échantillon de 14 galaxies à décalage spectral intermédiaire ($z \sim 0.7$) observées avec la l'instrument ACS/HST dans le champ profond GOODS-South, et classées comme granulaires. Chaque image est une composition réalisée à partir des filtres B, V et i . Crédits : Bournaud et al. (2012).

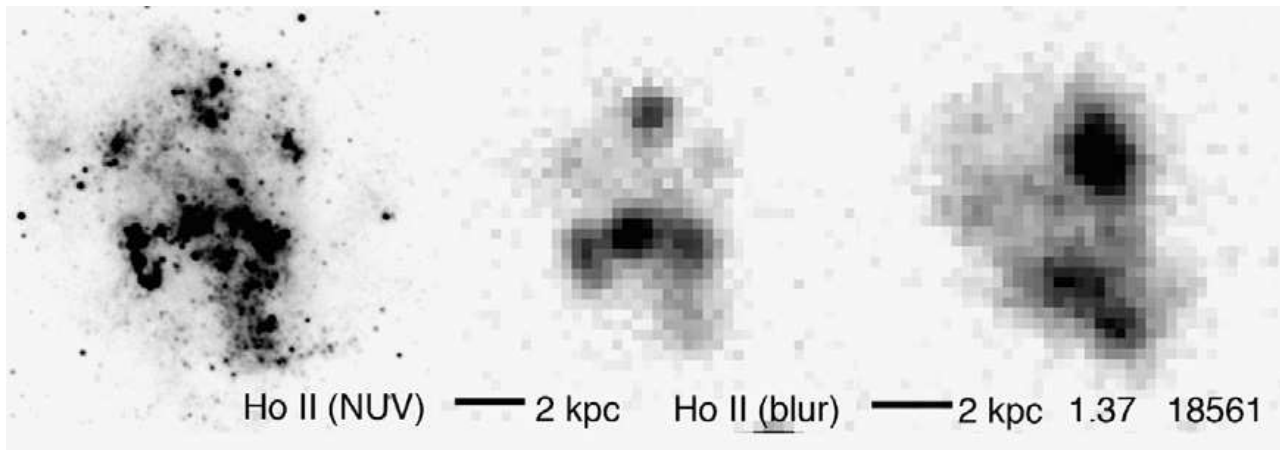


Figure 4.3 – Comparaison d’une galaxie naine irrégulière de l’Univers local à une galaxie granulaire à grand décalage spectral. À gauche : image dans l’UV proche de la galaxie naine irrégulière Ho II. À droite : amas granulaire 18561 du programme GOODS vu dans la bande passante V606. Au milieu : image de la galaxie naine Ho II avec une résolution dégradée afin de pouvoir comparer à l’amas granulaire 18561. La résolution spatiale (780 parsecs) et la longueur d’onde dans le référentiel au repos ($\sim 2400 \text{ \AA}$) sont identiques à celles de l’image ACS de la galaxie 18561. Les asymétries morphologiques observées dans la galaxie naine locale sont comparables à celles de la galaxie en amas granulaire. Cependant, les galaxies locales irrégulières ont une masse plus faible d’un facteur 10-100. Crédits : Elmegreen et al. (2009b).

dans les densités de colonne. De telles vitesses turbulentes sont en accord avec les observations (Genzel et al., 2011 ; Epinat et al., 2012), et les densités de colonne observées dans les galaxies granulaires sont cohérentes avec celles observées dans les régions internes des disques de l’Univers local.

Le nombre croissant des observations spectroscopiques 3D de galaxies à grand décalage spectral a révélé une grande fraction d’objets montrant des champs de vitesses typiques d’une rotation ordonnée (figure 4.4). La présence d’une rotation ordonnée malgré la présence claire de granules dans les morphologies stellaire et du gaz ionisé renforce l’idée que ce type d’objets n’est pas le résultat d’une fusion récente, mais bien celui d’un processus d’évolution séculaire des disques, capables de se fragmenter tout en maintenant une rotation très similaire à celle des galaxies spirales. L’utilisation d’optique adaptative permet notamment aux spectrographes intégral de champs d’atteindre des résolutions spatiales rendant possible une analyse des spectres intégrés de chacun de ces granules. Genzel et al. (2011) observent dans quatre galaxies massives de l’échantillon SINS une composante large des raies d’émission $H\alpha$ et [NII] dont le meilleur ajustement est obtenu pour une Gaussienne d’une largeur de 500 km.s^{-1} , ce qui serait une indication de l’existence d’importants mécanismes d’éjection du gaz du fait d’une activité de formation stellaire intense dans les granules.

4.2 Formation des galaxies granulaires

Les simulations numériques sont un outil de choix pour comprendre la nature des processus physiques capables de fragmenter les disques jeunes et gazeux en de multiples granules.

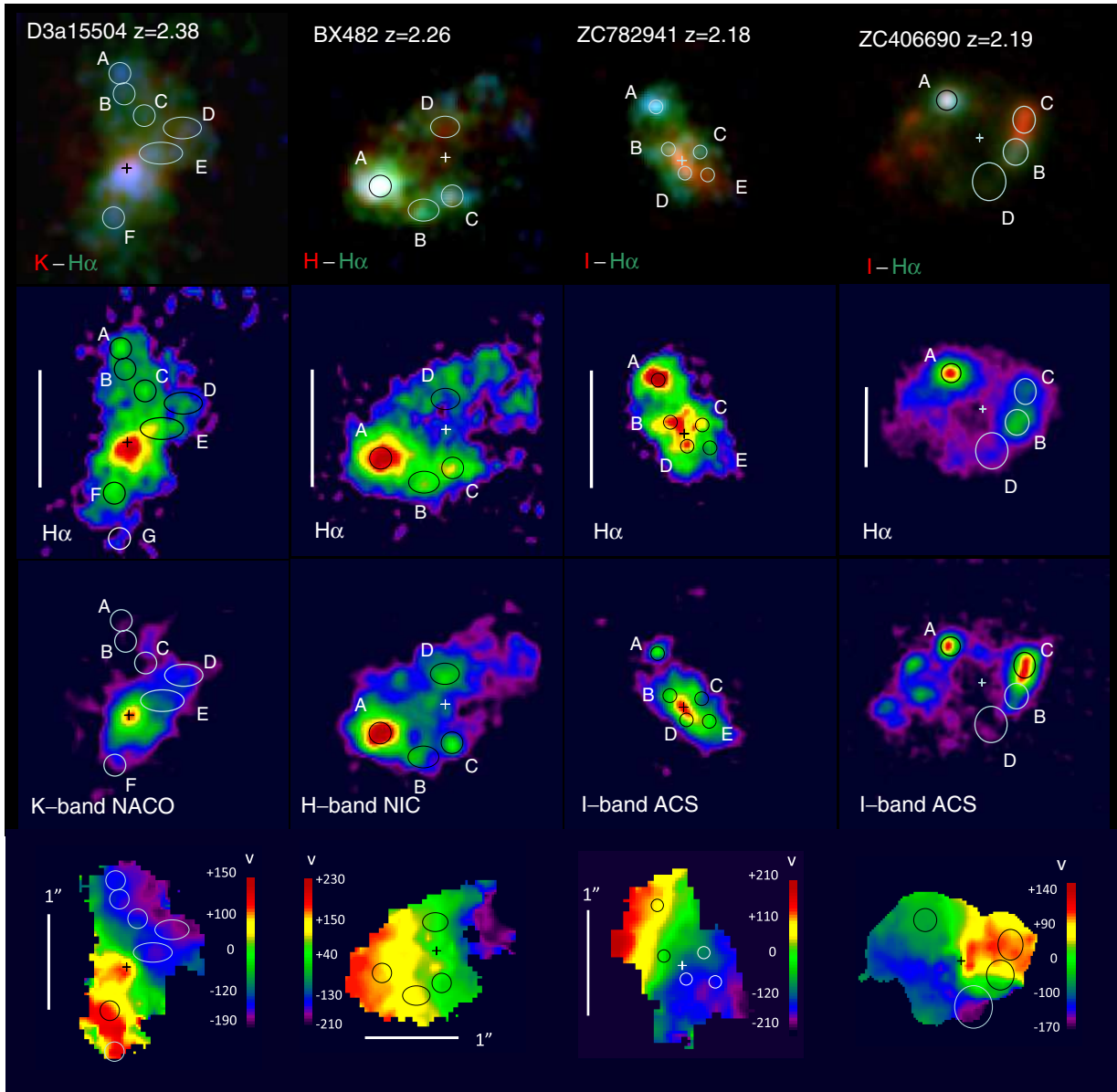


Figure 4.4 – Morphologie granulaire du gaz ionisé dans quatre galaxies de l'échantillon SINS. Ligne du haut : images composite $H\alpha$ et UV/optique du continu stellaire (dans le référentiel au repos) pour quatre galaxies massives avec un fort taux de formation stellaire à $z \sim 2$. Seconde ligne : Flux intégré $H\alpha$ observé avec l'instrument SINFONI. Troisième ligne : images du flux du continu stellaire UV/optique avec le même code de couleur que les images du flux $H\alpha$. Ligne du bas : champs de vitesses des galaxies, obtenus par l'ajustement Gaussien sur la raie $H\alpha$. Les granules les plus massifs sont identifiés par les cercles annotés des lettres A à F. Les différentes bandes passantes sont indiquées sur chaque image. La croix sur chacune des images indique le centre cinématique de la galaxie. Crédits : Genzel et al. (2011).

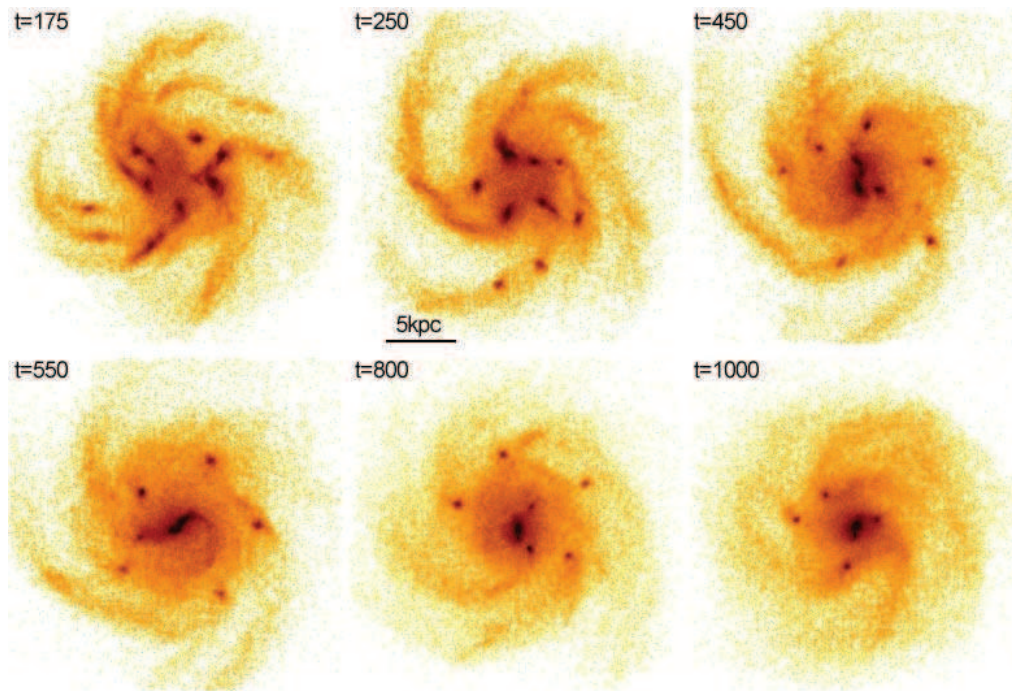


Figure 4.5 – Projections de face de la densité de masse (étoiles et gaz) pour une simulation modélisant l'évolution d'un disque (gaz et étoiles) avec une masse de $7 \times 10^{10} M_{\odot}$. Les granules se forment rapidement, et se déplacent vers le centre de la galaxie, où ils fusionnent et forment un bulbe en moins de 1 Gyr. Quelques granules demeurent dans le disque à la fin de la simulation. Crédits : Elmegreen et al. (2008).

La fragmentation du disque est en réalité un processus naturel relativement simple à obtenir dans une simulation du moment que l'on dispose d'une résolution numérique permettant de résoudre la structure verticale du disque, et que l'on permette au gaz de dissiper son énergie interne par refroidissement radiatif (Bournaud et al., 2007a). En effet, les disques de l'Univers lointain possèdent des fractions de gaz élevées (Daddi et al., 2010a) et sont situés d'après le modèle cosmologique Λ -CDM dans des halos de matière sombre bien moins massifs que ceux des galaxies locales. Le gaz étant une composante radiative, il est capable de se refroidir et donc de subir un effondrement gravitationnel. Cet effondrement gravitationnel fragmente le disque en plusieurs granules, aussi bien pour la composante stellaire que gazeuse, qui poursuivent leur rotation autour du centre de masse de la galaxie. Ces granules sont capables d'échanger du moment angulaire avec le disque diffus et de migrer progressivement vers le centre de la galaxie (figure 4.5). Pour que ces instabilités se développent, il est nécessaire d'avoir un paramètre de Toomre (chapitre 2) $Q < 1$. L'existence d'une distribution de masse sous forme sphéroïdale (halo stellaire, halo de matière sombre) a tendance à stabiliser le disque contre les instabilités axisymétriques qui tendent à le disloquer. A contrario, une densité de masse surfacique élevée dans un plan a tendance à rendre le disque instable. La stabilisation artificielle de la composante gazeuse réalisée en bloquant numériquement le refroidissement empêche la formation des granules et produit des galaxies beaucoup plus similaires morphologiquement à des spirales (figure 4.6). On peut définir la longueur de Jeans pour une instabilité gravitationnelle :

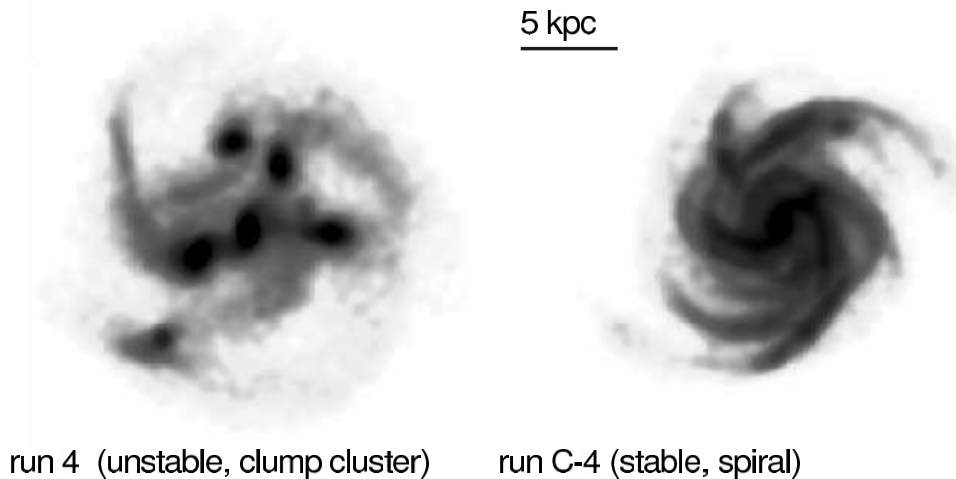
Morphology at $t=0.5$ Gyr

Figure 4.6 – Deux cartes de densité de masse, pour un modèle numérique de galaxie avec une masse initiale de $7 \times 10^{10} M_{\odot}$ et une échelle de longueur initiale de 6 kpc. À gauche : modèle numérique avec refroidissement. À droite : modèle numérique équivalent mais stabilisé. Les deux projections sont réalisées toutes les deux 500 Myr après les conditions initiales. Crédits : Bournaud et al. (2007a).

$$\lambda_J = \frac{\sigma^2}{\pi G \Sigma}, \quad (4.1)$$

avec σ la dispersion locale de vitesses, et Σ la densité surfacique locale. Pour une galaxie avec une fraction de gaz $f_g \sim 50\%$, un rayon $R \sim 9$ kpc, une masse de gaz $M_g = 10 \times 10^{10} M_{\odot}$, et en supposant une distribution de masse uniforme ($\Sigma \sim M_g / (\pi R^2)$), on obtient $\sigma \simeq 50$ km.s $^{-1}$ en supposant $\lambda_J = 500$ pc. Ce calcul est cohérent avec les mesures de dispersions de vitesses réalisées dans les échantillons spectroscopiques à grand z (Förster Schreiber et al., 2009 ; Epinat et al., 2012), et suggère donc que le scénario de fragmentation par instabilités gravitationnelles serait à l'oeuvre dans les galaxies granulaires.

4.3 Propriétés des granules dans les simulations de disques isolés MIRAGE

Les conditions initiales des disques de l'échantillon MIRAGE ont été créées afin de pouvoir observer le comportement du gaz dans des galaxies granulaires isolées et en interaction. En effet, des galaxies avec une fraction de gaz initiale élevée, qui accrètent continuellement du gaz primordial du milieu inter-galactique et situées dans des halos de matière sombre peu concentrés ($\sim 30\%$ de la masse présente dans le rayon optique) sont propices à l'apparition de granules formés par effondrement gravitationnel. La question qui se pose est de savoir si ces granules sont détruits au-delà de leur temps dynamique (~ 50 Myr), de manière similaire aux nuages moléculaires des galaxies spirales de l'Univers local, où bien si ils sont en mesure de survivre aux mécanismes de rétro-action stellaire durant plus d'un temps orbital du disque (~ 300 Myr). Dans ce dernier cas de figure, les granules seraient capables de migrer

vers le centre de masse du disque et contribueraient à la construction du bulbe stellaire central. Dans la publication suivante, nous fournissons des éléments de réponse en procédant à l'analyse des simulations numériques de disques isolés de l'échantillon MIRAGE, présentées dans le chapitre précédent. Notamment, l'utilisation d'un modèle physique de pression radiative entretenue par les étoiles jeunes de type OB permet de poser un regard nouveau sur cette problématique récente. Dans les simulations MIRAGE, nous trouvons une expulsion de gaz continue comparable aux taux de formation stellaire dans les granules, avec des vitesses permettant au gaz de s'échapper de la galaxie. Les granules perdent aussi de la masse par interaction gravitationnelle et les populations stellaires contenues dans les granules restent relativement jeunes (≤ 200 Myr), comme cela est suggéré par les observations Wuyts et al. (2012). Les granules survivent à la déplétion en gaz du fait qu'ils sont en orbite dans un milieu turbulent constamment ré-enrichi par l'accrétion de gaz chaud du milieu inter-galactique. Ce milieu environnant permet une compensation des pertes de masse dues aux injections d'énergie successives des étoiles jeunes. Les simulations de disques isolés de l'échantillon MIRAGE produisent des taux d'éjection de gaz avec des propriétés comparables aux rares données disponibles dans les observations, tout en suggérant que les granules peuvent survivre plusieurs centaines de millions d'années et terminer leur migration vers le centre de la galaxie.

Ma contribution à cette publication a essentiellement consisté à préparer les simulations numériques de l'échantillon MIRAGE, particulièrement adaptées à ce type d'analyse. Le choix des conditions initiales très riches en gaz, la résolution numérique élevée (~ 7 pc) ainsi que la modélisation de multiples processus de rétro-action stellaire en font un échantillon unique pour répondre aux questions que pose l'évolution des granules dans un disque avec un contexte d'accrétion de gaz chaud du milieu circum-galactique.

THE LONG LIVES OF GIANT CLUMPS AND THE BIRTH OF OUTFLOWS IN
GAS-RICH GALAXIES AT HIGH REDSHIFTFRÉDÉRIC BOURNAUD¹, VALENTIN PERRET², FLORENT RENAUD¹, AVISHAI DEKEL³, BRUCE G. ELMEGREEN⁴,
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ABSTRACT

Star-forming disk galaxies at high redshift are often subject to violent disk instability, characterized by giant clumps whose fate is yet to be understood. The main question is whether the clumps disrupt within their dynamical timescale (≤ 50 Myr), like the molecular clouds in today's galaxies, or whether they survive stellar feedback for more than a disk orbital time (≈ 300 Myr) in which case they can migrate inward and help building the central bulge. We present 3.5–7 pc resolution adaptive mesh refinement simulations of high-redshift disks including photoionization, radiation pressure, and supernovae feedback. Our modeling of radiation pressure determines the mass loading and initial velocity of winds from basic physical principles. We find that the giant clumps produce steady outflow rates comparable to and sometimes somewhat larger than their star formation rate, with velocities largely sufficient to escape the galaxy. The clumps also lose mass, especially old stars, by tidal stripping, and the stellar populations contained in the clumps hence remain relatively young (≤ 200 Myr), as observed. The clumps survive gaseous outflows and stellar loss, because they are wandering in gas-rich turbulent disks from which they can reaccrete gas at high rates compensating for outflows and tidal stripping, overall keeping realistic and self-regulated gaseous and stellar masses. The outflow and accretion rates have specific timescales of a few 10^8 yr, as opposed to rapid and repeated dispersion and reformation of clumps. Our simulations produce gaseous outflows with velocities, densities, and mass loading consistent with observations, and at the same time suggest that the giant clumps survive for hundreds of Myr and complete their migration to the center of high-redshift galaxies. These long-lived clumps are gas-dominated and contain a moderate mass fraction of stars; they drive inside-out disk evolution, thickening, spheroid growth, and fueling of the central black hole.

Key words: galaxies: bulges – galaxies: evolution – galaxies: formation – galaxies: high-redshift – galaxies: structure

Online-only material: color figures

1. INTRODUCTION

Massive high-redshift galaxies often have irregular morphologies dominated by giant clumps in optical imaging surveys (Cowie et al. 1996; Elmegreen et al. 2007), ionized gas observations (Genzel et al. 2008), and resolved molecular gas maps (Tacconi et al. 2013). Typically, a clumpy star-forming galaxy of stellar mass of a few $10^{10-11} M_{\odot}$ at redshift $z \approx 1-3$ contains a handful of kpc-sized clumps containing each a few $10^{8-9} M_{\odot}$ of gas and stars (Elmegreen & Elmegreen 2005; Genzel et al. 2011). About half of the star formation occurs in the giant clumps, the other half being in an apparently diffuse disk probably consisting of smaller, unresolved gas clouds (Elmegreen et al. 2009). At last half of star-forming galaxies at $z > 1$ in the CANDELS near-infrared imaging survey are clumpy (M. Mozena et al., in preparation; Y. Guo et al., in preparation).

From the following set of evidence, it is now widely accepted that the majority of the clumps form in situ by gravitational instability with high characteristic Jeans mass and length (around $10^9 M_{\odot}$ and 1 kpc, respectively). The host galaxies apparent shapes range from round to linear, with the respective fractions expected for face-on to edge-on disks (with ≈ 1 kpc scaleheights; Elmegreen et al. 2004) and disk-like kinematics with high

turbulent velocity dispersions in both H α and CO (Genzel et al. 2006; Shapiro et al. 2008; Bournaud et al. 2008; Contini et al. 2012; Daddi et al. 2010a; Epinat et al. 2012; Tacconi et al. 2013, with perhaps some redshift evolution, Puech 2010). The high fractions of dense molecular gas (about half of the baryonic mass; Daddi et al. 2008, 2010b; Tacconi et al. 2010) naturally explain the occurrence of these violent instabilities and the high associated velocity dispersions required to regulate the disk at a Toomre stability parameter $Q \approx 1$ (Dekel et al. 2009a, 2009b; Bournaud & Elmegreen 2009; Elmegreen & Burkert 2010; Burkert et al. 2010) and observations are consistent with a dynamical state of the gas close to the $Q \approx 1$ instability limit (Genzel et al. 2011). Simulations with different numerical schemes, of both idealized galaxies and galaxies in cosmological context, show the rapid development of strong turbulence and giant clump formation in such gas-rich galaxies (Noguchi 1999; Bournaud et al. 2007; Agertz et al. 2009; Ceverino et al. 2010; Powell et al. 2011; Martig et al. 2012; Dubois et al. 2013; Hopkins et al. 2012).

The violent clump instability thickens the stellar disk (Bournaud et al. 2009) and drives an inflow toward the nucleus at a much higher rate than spiral arms and bars (Krumholz & Burkert 2010; Bournaud et al. 2011). However, other major dynamical effects depend on the lifetime of the giant clumps

and their ability to survive stellar feedback. Early simulations with no or weak feedback found that the clumps retain much of their mass or increase their mass over time, and survive feedback (Immeli et al. 2004; Elmegreen et al. 2008a; N. Mandelker et al., in preparation). In this case, the clumps migrate radially through dynamical friction and gravitational torques in only a few 10^8 yr and convey important amounts of gas and stars inward. This increases the general instability-driven inflow; can rapidly (≤ 1 Gyr) reshape the disk radial profile into an exponential (Bournaud et al. 2007); form a massive classical, low-Sersic-index bulge (Noguchi 1999; Elmegreen et al. 2008a; Ceverino et al. 2010); and increase the nuclear gas flows, powering bright active galactic nucleus (AGN) episodes (Gabor & Bournaud 2013; Dubois et al. 2012).

However, the ability of giant clumps to survive stellar feedback for 10^8 yr or more has been questioned. The momentum conveyed from young stars by photons is an efficient process to affect dense gas structures even before supernovae explode (Murray et al. 2010), and this has been suggested to rapidly disrupt the giant clumps of high-redshift galaxies where high gas densities could trap the photons (Murray et al. 2011). A potential issue with short-lived clump scenarios is that observed giant clumps need to live long enough to accumulate their high baryonic masses. Simulations with strong ad hoc radiative feedback have shown that giant clumps could form and disrupt within 100 Myr (Genel et al. 2012b; Hopkins et al. 2012), but leaving unclear whether rapidly disrupted clumps could reform sufficiently quickly to explain the high observed fraction of clumpy galaxies.

Two observations have raised the possibility of a short clump lifetime. First, outflows have been detected in the vicinity of a few massive clumps, with outflow rates of the order of their star formation rate (SFR) or a few times higher (Genzel et al. 2011; Newman et al. 2012a, 2012b), and with high electronic temperatures suggesting strong pressure gradients (Lehnert et al. 2013). Second, the clumps are much fainter in near-infrared imaging (Elmegreen et al. 2009), indicating that their masses are relatively limited in the old stellar components, and the typical ages of the stellar populations contained in clumps are rarely larger than ≤ 200 Myr (Wuyts et al. 2012). This indicates that the clumps do not disrupt as rapidly as in simulations with extreme ad hoc feedback (Genel et al. 2012a) and may be rather consistent with typical clump migration timescales of 200–400 Myr (Bournaud et al. 2007; Dekel et al. 2009b; Ceverino et al. 2010).

In contrast, it has been proposed that a massive clump wandering in a turbulent disk may rapidly accrete significant amounts of gas (Dekel & Krumholz 2013), which may compensate for the losses in star formation and outflows and make clumps longer lived. Even if initially warm, gas (re)accreted this way should become dense enough to cool in less than a dynamical time, hence rapidly refueling the bound star-forming component in the clumps (Pflamm-Altenburg & Kroupa 2009).

In this paper, we present high-resolution adaptive mesh refinement (AMR) simulations of high-redshift clumpy galaxies using a new, physically motivated modeling of stellar feedback (Section 2). We show that giant clumps produce outflows and lose aged stars by tidal stripping, but keep their mass and SFR about constant over time owing to continuous reaccretion of gas from the surrounding disk (Section 3). The clumps in these models are long-lived and complete their migration inward until they merge with other clumps or with the bulge. Also, they have properties (outflow rates, SFRs, stellar populations,

contrast in optical imaged and stellar mass maps) that appear fully consistent with observations (Section 4).

2. SIMULATIONS

2.1. Models of High-redshift Disks

In this study, we follow the detailed evolution of giant clumps in high-redshift galaxies, in particular their response to star formation and feedback processes. To this aim we use “idealized” (closed-box) simulations of systems with global sizes, masses, and gas fractions similar to real galaxies at redshift two, which naturally form giant clumps similar to the observed ones as studied by many previous authors. The advantage of such idealized simulations is that the interstellar medium (ISM) structure can be resolved with a resolution a decade finer than in state-of-the-art cosmological simulations, resolving the size of heated structures around young star clusters, and sub-structures in the giant clumps through which outflows can propagate. These simulations model disk galaxies with mass distributions, rotation curves, and gas fractions that are realistic for $z = 1$ – 3 galaxies, although their formation history at higher redshifts is not followed. During the evolution of their giant clumps, these simulations lack the replenishment of the disk by external gas inflow, which may limit the ability of the clumps to reaccrete gas from the surrounding regions of the disk, thus providing a lower limit to the survivability of clumps.⁷ As for dynamical aspects, various studies have shown that both the gravity and turbulence of these galaxies are controlled by internal physics rather than by external inflow (e.g., self-gravity: Elmegreen & Burkert 2010, feedback: Hopkins et al. 2013; Genel et al. 2012a). The disk galaxy dynamics can therefore be followed quite reliably in isolation for several hundred Myr.

The fiducial high-redshift disk simulations studied here, labeled G1, G2, and G3 and displayed in Figures 1–3, are taken from the MIRAGE sample of disk and mergers presented in Perret et al. (2013). An additional model G’2, close to G2 but using a different combination of feedback and resolution parameters, is also used here.

The initial baryonic masses of galaxies G1–G3 are 8.7, 3.5, and $1.4 \times 10^{10} M_{\odot}$, respectively. Their evolution starts with a gas fraction (gas/gas+stars) of 60%. The disk scale lengths are set according to the Vergani et al. (2012) scaling relations, equally compatible with other high-redshift scaling relations (e.g., Dutton et al. 2011). The bulge-to-disk stellar mass ratio is 30%. The star formation model subsequently detailed naturally leads to a realistic specific SFR of the order of one Gyr^{-1} , realistic for the main sequence at redshift $z \approx 1$ – 2 (Daddi et al. 2007; Elbaz et al. 2011; Nordon et al. 2013). The evolution is followed for 1 Gyr, and the average gas fraction over this period is 47%. Stars and dark matter are modeled with about 1 million particles each (see details in Perret et al. 2013).

The alternate model G’2 starts with initial conditions similar to G2 except that the total mass is 17% higher and the initial gas fraction is 50%; the main difference is the (weaker) supernovae feedback modeling as subsequently detailed. It also uses a twice higher spatial resolution.

The model parameters are summarized in Table 1.

⁷ This effect should not be major because the gas fraction decreases slowly compared with the 400–500 Myr timescale required for clump migration. In cosmological simulations, the inability to explain the very high gas fractions of $\approx 50\%$ at $z \approx 2$ because of too early consumption into star formation at $z > 3$ may be a larger issue regarding the galactic reservoir of gas in which giant clumps form and evolve (e.g., Ceverino et al. 2010).

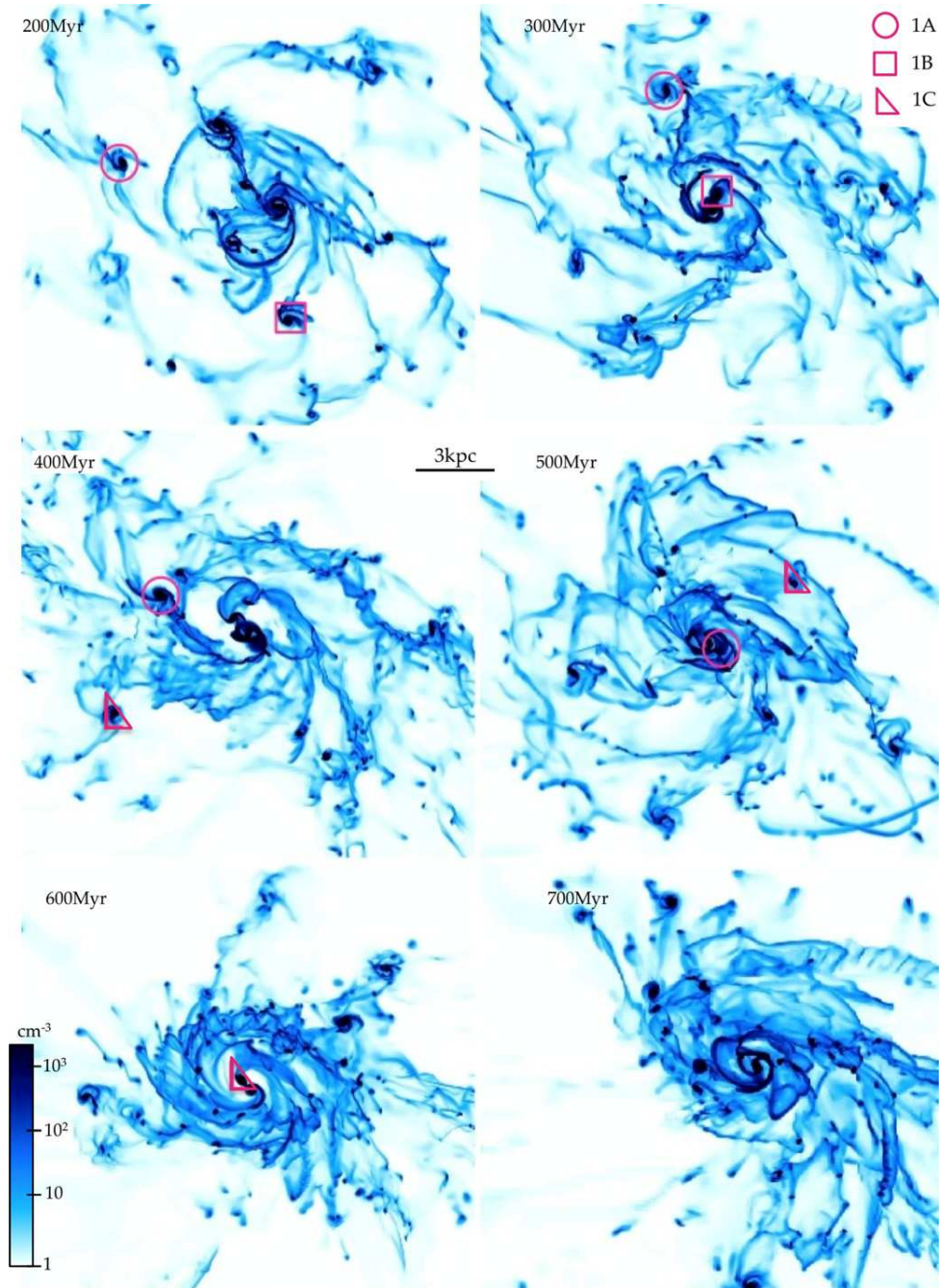


Figure 1. Sequence of snapshots showing the mass-weighted average of the gas density along each line of sight (i.e., the typical three-dimensional density reach along each line of sight, rather than the column density) for model G1 (high mass). Snapshots are taken every 100 Myr from 200 to 700 Myr after the beginning of the simulation. Three typical clumps were selected for analysis, marked with symbols on the maps. At $t = 700$ Myr, these three typical clumps have merged with the central disk or bulge or with other giant clumps, hence being unmarked in the last snapshot, on which other clumps formed later on in the outer disk material and/or recycled gas are still present. Detailed sequences and movies of our fiducial models are available in Perret et al. (2013) and allow the reader to track long-lived clumps in detail.

(A color version of this figure is available in the online journal.)

2.2. AMR Simulation Technique

We use the AMR code RAMSES (Teyssier 2002) with a quasi-Lagrangian AMR refinement methodology similar to that used for idealized galaxy simulations in Teyssier et al. (2010), Bournaud et al. (2010), and Renaud et al. (2013). The grid resolution reaches 7 pc in the high-density and/or Jeans-unstable regions for our fiducial models G1 to G3. Model G'2

uses a 3.5 pc resolution. Starting from the coarse level, each AMR cell is refined into 2^3 new cells if (1) it contains more than 25 particles, (2) its gas mass content is larger than the “gas mass resolution” set equal to $1.5 \times 10^4 M_{\odot}$, or (3) the local thermal Jeans length is smaller than four times the current cell size. We also impose a pressure floor to keep the Jeans length greater than four times the smallest cell size and avoid spurious instabilities (Truelove et al. 1997, see Section 2.1 in Teyssier et al. 2010).

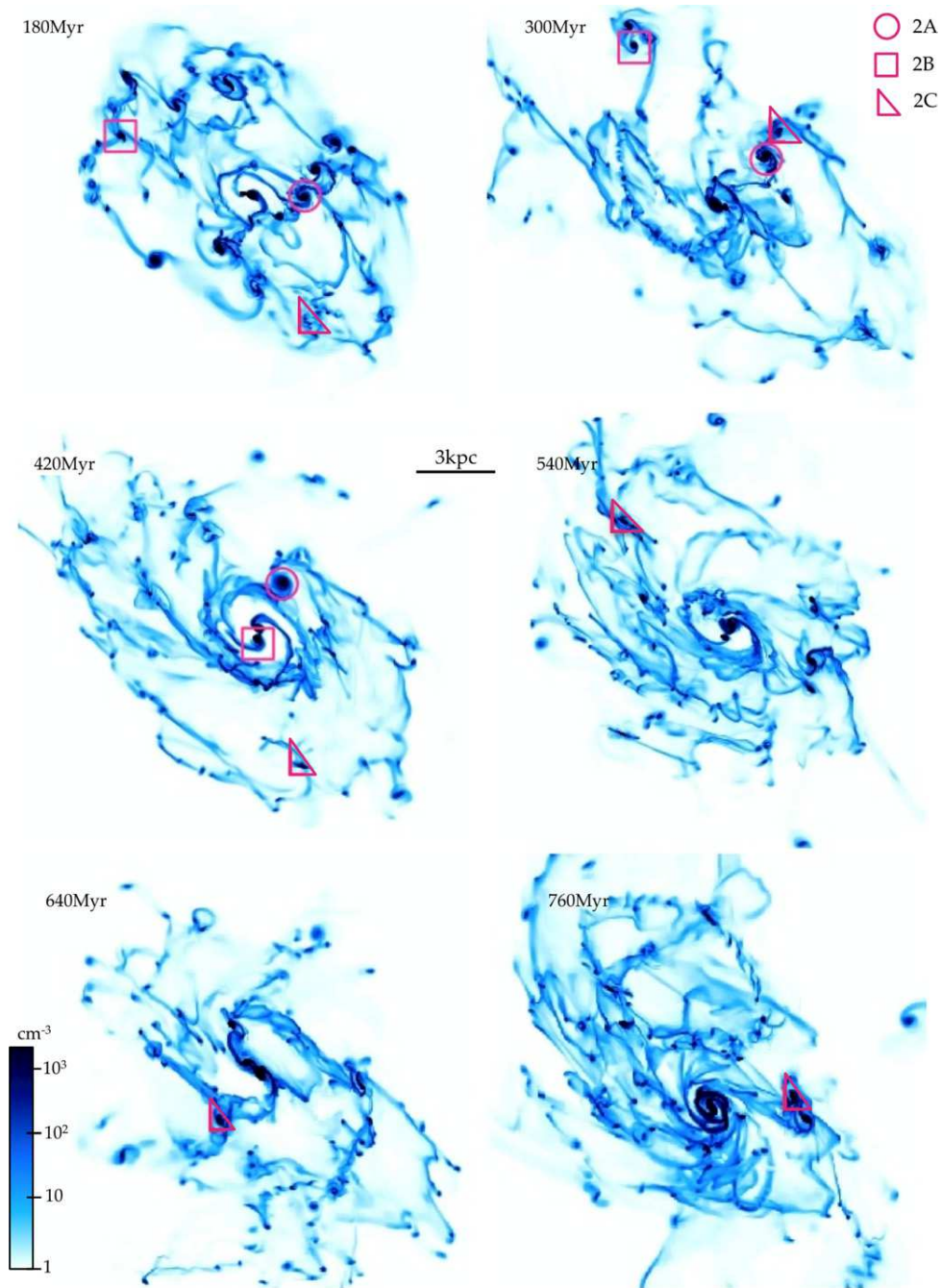


Figure 2. Same as Figure 1 for galaxy G2 (medium mass). Detailed sequences and movies of our fiducial models are available in Perret et al. (2013). (A color version of this figure is available in the online journal.)

Table 1
Model Parameters

Model	Stellar Mass ($10^9 M_{\odot}$)	Gas Mass ($10^9 M_{\odot}$)	Bulge Fraction (in Stellar Mass)	Supernova Feedback	Resolution (pc)
G1	87	130	7.5%	Reduced cooling	7
G2	35	52	7.5%	Reduced cooling	7
G'2	44	44	13%	Thermal cooling	3.5
G3	14	21	7.5%	Reduced cooling	7

Notes. Initial mass in stars and gas, bulge fraction (in the stellar mass distribution), cooling scheme for regions affected by supernovae explosions: unmodified thermal feedback, or reduced feedback to model non-thermal dissipation (Teyssier et al. 2013), maximal AMR resolution.

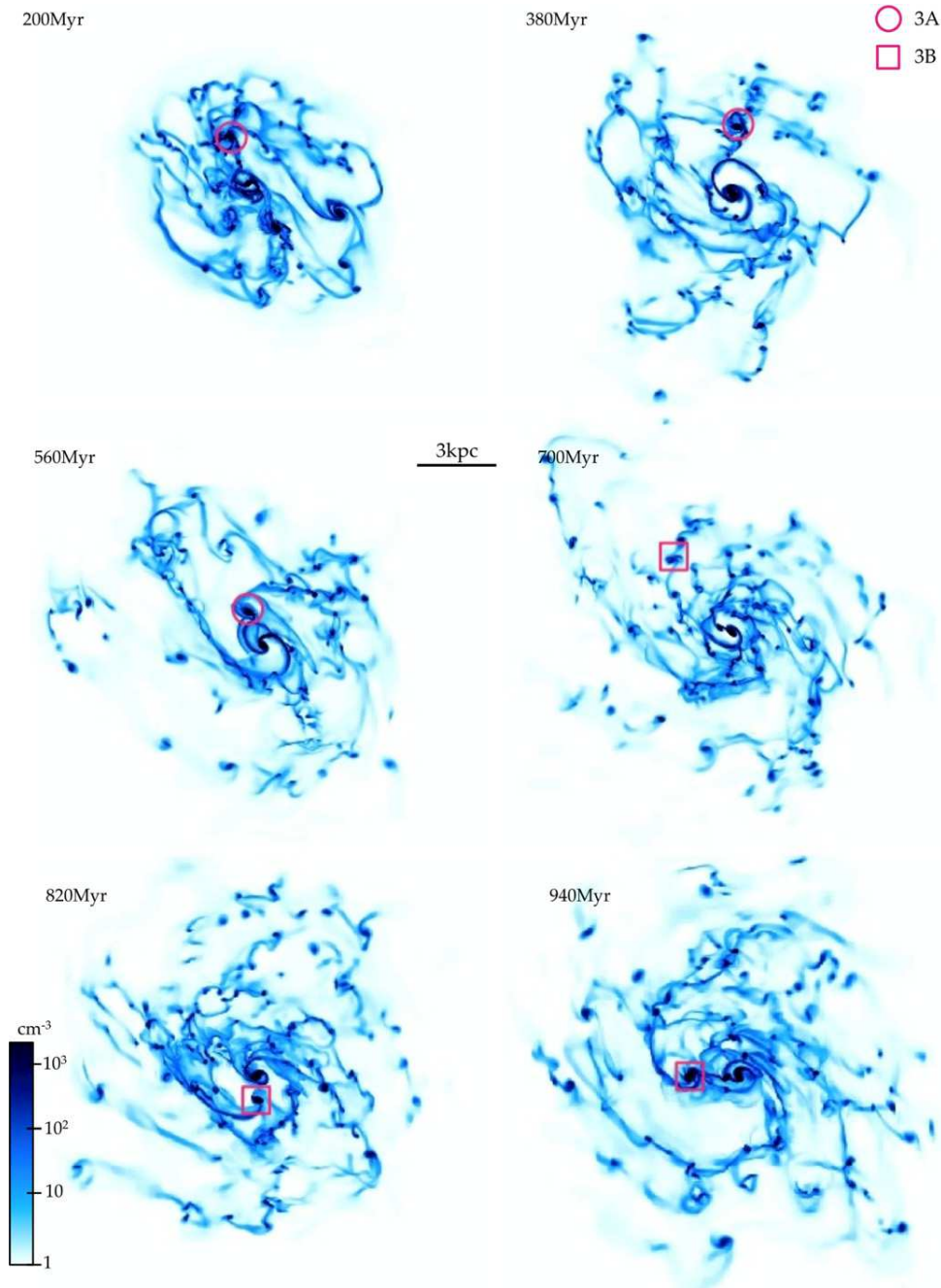


Figure 3. Same as Figure 1 for galaxy G3 (low mass). Detailed sequences and movies of our fiducial models are available in Perret et al. (2013).
(A color version of this figure is available in the online journal.)

We compute thermal evolution including atomic and fine-structure cooling assuming solar metallicity⁸ Such modeling naturally leads to a multiphase ISM, in which structures with number densities up to 10^{5-6} cm^{-3} are resolved. The initial gas in the disk has number densities of a few cm^{-3} so that we set the initial temperature to 10^4 K , which is the typical equilibrium temperature for such moderate densities (e.g., Bournaud et al. 2010, a lower temperature phase appears mostly once denser

structures arise). We prevent gas cooling below 100 K for better computational efficiency.⁹

2.3. Star Formation and Feedback

Star formation is modeled using a constant fraction $\epsilon_{\text{SF}} = 1\%$ of gas is converted into stars per local gravitational freefall time (2% in the alternative model G'2). This type of local model explains the observed relations between the surface densities of gas and SFR in a large range of objects (e.g., Elmegreen 2002;

⁸ The metallicity of massive galaxies at redshift 1–2 is high enough for their cooling not to be strongly reduced compared with solar metallicity (e.g., Erb et al. 2006).

⁹ Resolved structures of a few parsecs at 1000 K are dominated by turbulent pressure rather than thermal pressure, so the details of cooling below 1000 K have little net effect at the scales resolved in our models.

Renaud et al. 2012; Krumholz et al. 2012), and when applied to the current simulations matches the gas consumption timescales observed for main sequence galaxies at redshift $z \simeq 1-2$ (see Perret et al. 2013 for the current sample).

Three stellar feedback mechanisms are included:

1. *H II region photoionization*: around each stellar particle younger than 10 Myr, the photoionized region is computed using the Strömgren sphere approximation implemented by Renaud et al. (2013), the spherical approximation being justified by the fact that each H II region radius is not much larger than the numerical resolution used here (although large non-spherical regions do form from the overlap of smaller spherical regions around clusters of young stellar particles).
2. *Radiation pressure* from young stars is implemented using the Renaud et al. (2013) model. Once the available momentum in photons is computed for a standard initial mass function, we use the fact that most of the momentum is carried by ionizing photons and hence is acquired by the gas that is ionized, i.e., gas that lies in the previously calculated H II regions. This allows us to determine, from physical considerations, how the available momentum $m \times v$ is distributed into a mass m of gas pushed at a velocity v : in our model, m is linked to the mass of gas that each stellar particle is able to ionize. This is an important difference compared with other models of high-redshift disks with momentum-driven feedback, where the wind velocity was set by the escape velocity from the clump (or galaxy) to derive the mass loading factor (Genel et al. 2012b; Hopkins et al. 2012). For the same amount of available momentum, our model is able to generate a higher-velocity wind comprising a lower mass of gas, if high-density gas captures most of the momentum available from young stars. We use a trapping parameter $\kappa = 5$ to account for multiple scattering: This is rather high but is realistic as it compensates for other sources of momentum that are not explicitly included, such as (proto)-stellar winds (Krumholz & Thompson 2012; Dekel & Krumholz 2013).
3. *Supernovae* are implemented following the Dubois & Teyssier (2008) method, taking into account non-thermal processes as recently proposed by Teyssier et al. (2013). To account for non-thermal processes, the dissipation of the injected energy is limited to a fixed period of 2 Myr, slower than thermal cooling. This 2 Myr value is the typical turbulence dissipation timescale expected for the sub-grid structures that our simulations are unable to resolve (see details in Perret et al. 2013). This is qualitatively equivalent to the “delayed cooling” approach proposed by other authors to model the blastwave phase of supernovae explosions (e.g., Stinson et al. 2009), but the dissipation timescale is lower here because of the higher resolution.

Our alternative model G2 uses a pure thermal injection for supernovae feedback without a reduced dissipation rate. Hence, the supernova feedback in this model is weaker than in the fiducial runs, while the photoionization and radiative pressure feedbacks are the same.

2.4. Selection and Measurement of Giant Clumps

The giant clumps formed in the simulations are tracked visually with an output frequency of 4 Myr, and are analyzed in detail every 40 Myr. We do not consider here the clumps that undergo mergers with other clumps of similar mass in order to

make sure that a main clump progenitor can be unambiguously identified.

For each snapshot the visually determined position of each clump is used as an initial guess for its mass center. We compute the mass center of the baryons included in a sphere of radius 200 pc around this initial guess, and we iterate the procedure around the newly determined position, until the position shifts by a distance smaller than the resolution of the simulations. Once the mass center of the clump is found this way, we determine the clump radius and mass with the following procedure. Using a radius R_1 initially set to 100 pc, we compute the mass in the sphere of radius R_1 , and check the mass in a sphere with a volume that is twice as large (i.e., radius $2^{1/3} R_1$). If the mass increase is larger than 30%, we consider that the clump material was not mostly encompassed by the initial boundary, and we iterate the process with an initial radius $R_1 + 100$ pc, until the mass converges. Convergence is never reached at 100 pc, and always reached at radius ≤ 500 pc, namely the clump radii range from 200 pc to 500 pc. These positions and radii are used for all of the following measurements: gaseous and stellar masses, mass flows, and SFR per clumps. Given the method used, clump masses are determined with an uncertainty better than 30%; the uncertainty on clump radii can be up to 100 pc but only the masses are relevant in the following study. As an alternative technique, we used the definition of clump radius and mass proposed by Ceverino et al. (2012). The clump masses estimated this way are on average 23% higher but follow a very similar time evolution, as shown in the next section.

3. RESULTS

3.1. Short-lived Gas Clouds and Long-lived Giant Clumps

In our simulations, gas clouds below a mass of a few $10^7 M_\odot$ are short-lived. This is also true for gas clouds in low-redshift galaxies with the same technique (e.g., Renaud et al. 2013). A detailed time sequence showing the rapid formation and dissolution of low-mass gas clouds (not giant clumps) in our current model G2 is shown on Figure 4. Note that simulated molecular clouds in low-redshift galaxies were found to be short-lived and unvirialized even when only some of the feedback modes are included (e.g., Tasker 2011; Bournaud et al. 2010; Ceverino et al. 2012).

In contrast, the giant clumps more massive than $\approx 10^8 M_\odot$ remain long-lived in our simulations (Figures 1–3). Such clumps form by gravitational instability in high-redshift disks, with high Jeans length and mass because of the high level of turbulence¹⁰ required to self-regulate the disk at a Toomre parameter $Q \approx 1$. The average value of the gas-velocity dispersion¹¹ at the disk scale length and after two rotation periods ranges from 38 km s^{-1} (G3) to 53 km s^{-1} (G1). This process of giant clump formation in a turbulent medium has been studied in many works and is not detailed again here (Noguchi 1999; Immeli et al. 2004; Bournaud et al. 2007; Agertz et al. 2009; Dekel et al. 2009b; Ceverino et al. 2010; Genel et al. 2012b; Hopkins et al. 2012).

These giant clumps are long-lived and persist until coalescence with the central bulge or with another bigger clump. Most of them can be tracked in the simulations for 200–500 Myr and sometimes up to 700 Myr. We display in Figure 5 a time series showing the detailed evolution of a very long-lived

¹⁰ Regardless of the exact source of turbulent energy, gravity and star formation are the main internal energy sources and saturate at a $Q \approx 1$ level.

¹¹ One-dimensional rms dispersion measured in boxes of $(200 \text{ pc})^3$.

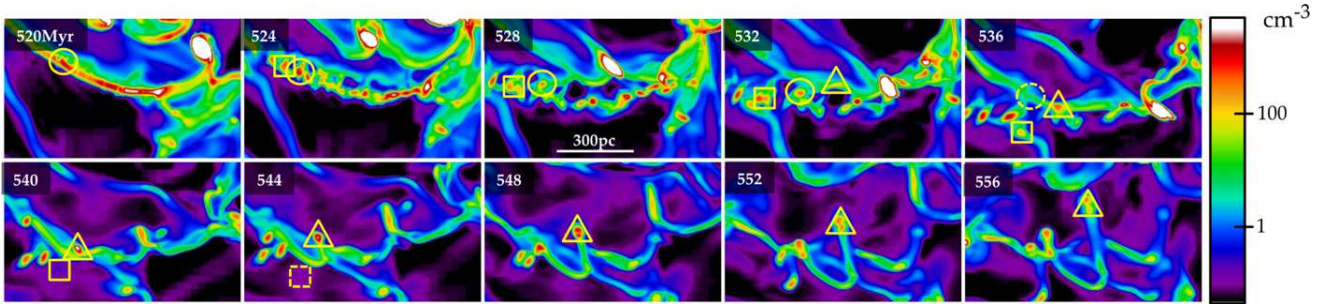


Figure 4. Sequence detailing the continuous formation and rapid dissolution of low-mass clouds, at the middle of the time evolution of model G2. The snapshot size is 0.8×1.6 kpc, and time is indicated in Myr, with one snapshot every 4 Myr. A moderately dense spiral arm, free of giant clump, forms a first generation of clouds. Two of these clouds are identified with the circular and boxy shapes: they leave only weak and dissolving knots after 15–20 Myr. Other clouds form at different locations, such as the one first indicated with the triangle at $t = 532$, which again dissolves, leaving a lower-density knot seen at $t = 556$ and unseen at $t = 600$ Myr.

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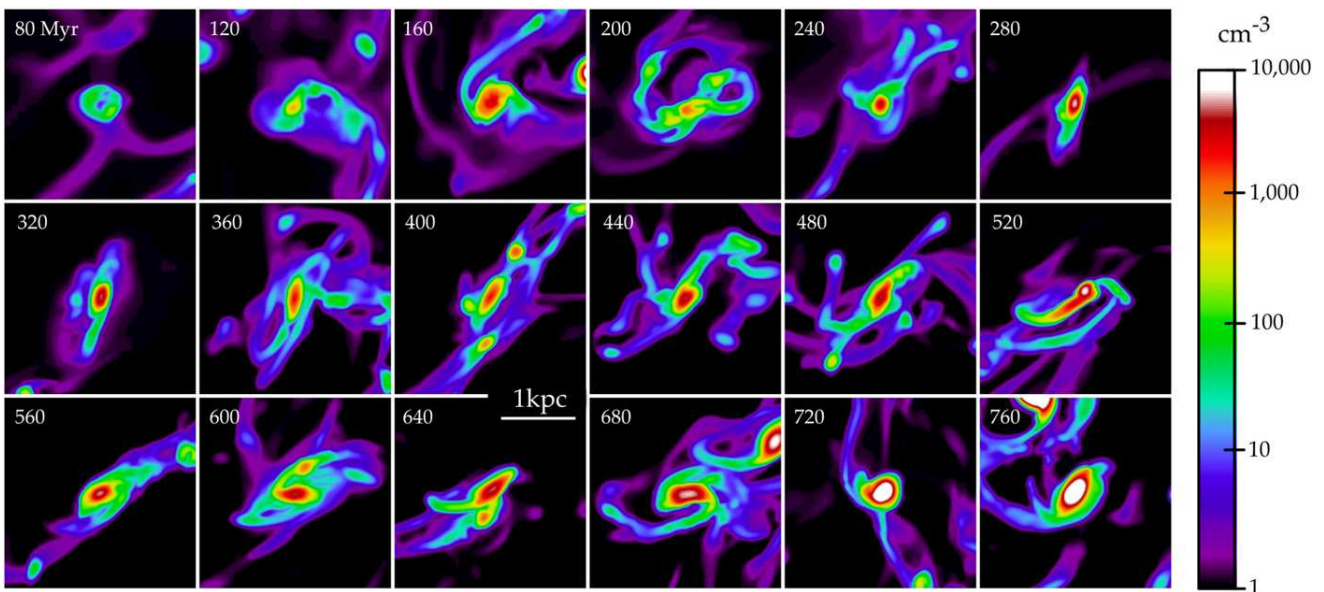


Figure 5. Zoomed views of gas in the long-lived clump 2C, with an average baryonic mass of $8 \times 10^8 M_{\odot}$. The snapshots show the mass-weighted average gas density, with one snapshot every 40 Myr. Between the third and fourth panels ($t = 160$ – 200 Myr), the clump accretes another clump (about half its mass), which triggers an increase in its SFR, and a later increase in the local outflow rate (see Figure 8); the clump gets a more disturbed appearance but the baryonic potential well in place rapidly reaccretes gas and the clump survives this local enhancement of the stellar feedback. Another such event, triggered by the accretion of surrounding diffuse gas and small clouds, occurs between the seventh and eighth panels ($t = 320$ – 360 Myr).

(A color version of this figure is available in the online journal.)

clump in our sample (clump 2C from model G2 displayed in Figure 2). The reason why this clump is particularly long-lived (at least 700 Myr) is that it forms at a large radius and keeps a low-eccentricity orbit, maximizing the timescale for inward migration by dynamical friction, while most giant clumps would reach the central bulge within 500 Myr.¹² As will appear in the following parts, this clump does not have extreme properties in terms of internal mass, size, or formation rate, and is simply used to illustrate the evolution of its mass content on a timescale that is not limited by rapid central coalescence, as would be the case with some clumps formed on different radii and/or on different orbits.

Over their long lifetime, the mass of the giant clumps remains relatively constant, without a major increase or decrease over time, in spite of the gas outflows detailed hereafter. We show in Figure 6 the time evolution of the baryonic mass of each of the clumps tracked in detail in our models: these masses, ranging from a few 10^8 to $2 \times 10^9 M_{\odot}$ and occasionally a bit

higher, and fluctuate about roughly constant values, without any significant global mass increase or decrease. As we detail in the following sections, giant clumps lose mass through high-velocity gas outflows and dynamical escape of aged stellar populations, but they also accrete gas from the surrounding disk. The rate of mass accretion in our simulations (detailed in Section 3.4) compensates for the losses and is consistent with the theoretical estimates from Dekel & Krumholz (2013) or Pflamm-Altenburg & Kroupa (2009). Hence, realistic feedback does not disrupt the giant clumps on timescales shorter than 10^8 yr, even when radiative pressure and non-thermal effects in supernova bursts are included, and continuous capture of baryons maintains their mass high on the long term.

3.2. Gas Outflows from Massive Clumps

The time evolution of the clump masses (Figure 6) shows that, despite being on average roughly constant, their baryonic mass can sometimes decrease. The detailed sequence of clump 2C (Figure 5) also shows that the mass and density of the gas

¹² For galaxies in the mass range studied here.

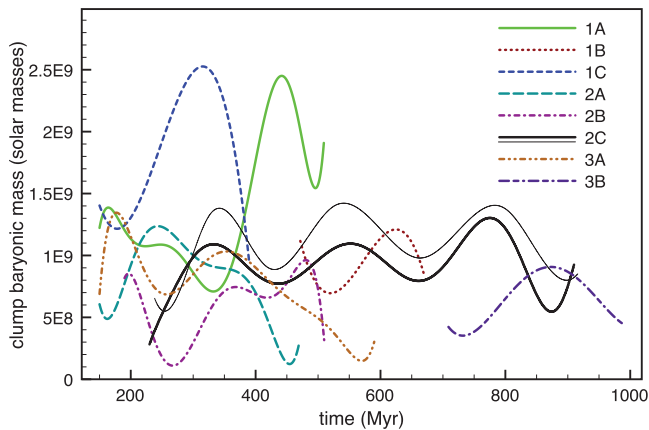


Figure 6. Evolution of the baryonic (gas+stars) mass of clumps as a function of time for the clumps tracked in the simulations. Measurements were averaged over bins of 15 Myr and polynomial interpolation was used. The thin line for clump 2C uses an alternative measurement of clump mass and radius (see text), yielding slightly higher clump masses but very similar evolution a mass fluctuating about the initial Jeans mass.

(A color version of this figure is available in the online journal.)

in the clumps can sometimes be significantly reduced. While the reduction of the gas mass alone could be the result of consumption by star formation, the fact that the total baryonic mass can sometime decrease implies that gas can actually be lost by the clumps, not just converted into stars.

The giant clumps in our simulations produce gaseous outflows. The birth and evolution of high-velocity outflows from the clumps is illustrated in Figure 7, where the same clump is shown after 90 Myr (when it just formed) and 270 Myr of evolution. At the early stage, shock fronts are formed by the outflowing gas onto the hot and diffuse halo surrounding the galaxy, with number densities of 10^{-2} – 10^{-1} cm^{-3} and velocities around 200 km s^{-1} . Later on, a higher velocity outflow escapes the clump region at velocities of 300 – 400 km s^{-1} with densities of a few 10^{-2} cm^{-3} , corresponding to an outflow rate of 1 – $2 M_{\odot} \text{ yr}^{-1}$ across a section of 1 kpc^2 . It is interesting to note that the gaseous outflow (1) can largely exceed the local escape velocity and (2) continues to be accelerated more than 1 kpc above the disk mid-plane, because of a pressure gradient that develops in the steady-state outflow regime. This turns into a galactic-scale outflow, the properties (rate, velocity) of which will be analyzed elsewhere (Perret et al. 2013). A movie is available¹³ that shows the formation of giant clumps followed by the development and expansion of gaseous outflows from clumps to galactic scales in model G².

We have systematically measured the outflow rate from each clump in two ways: through a spherical boundary around the clump (identical to the boundary used to measure the clump mass) and through a pair of $1 \text{ kpc} \times 1 \text{ kpc}$ planar boundaries positioned¹⁴ 1 kpc above and below the disk mid-plane. In the following sections, we use the latter measurement, after noticing that the former definition leads to similar measurements of the outflow rates from clumps with some additional uncertainty¹⁵

(on average 23% higher, with an rms relative deviation between the two measurements of 37%). The fact that these measurements yield similar results indicates that the outflows are significantly non-isotropic even at only 1 kpc from the clump centers, as seen in the examples of outflow velocity fields displayed in Figure 7.

The typical outflow rate from the giant clumps in our simulations is 1 – $4 M_{\odot} \text{ yr}^{-1}$, and up to $12 M_{\odot} \text{ yr}^{-1}$ during short episodes, of the order of the SFR in each clump and up to a few times higher. The time evolution of the outflow rate is shown for three individual clumps on Figure 8. Figure 9 shows the statistical distribution of the outflow rate to SFR ratio for our main sample and for simulation G² using weaker supernovae feedback (and higher resolution).

We show in Figure 10 the mass-weighted point distribution function of the gas velocity along a line of sight crossing the clump center, using a circular aperture of 300 pc radius around the line of sight, and with a line of sight inclined by 20° from the disk axis. This is equivalent to a pseudo-observed “spectrum” of the clump with a quasi-face-on orientation of the galaxy and a 600 pc FWHM beam,¹⁶ using crude linear conversion of gas mass into flux. For a typical, quiescent phase, the spectrum shows a narrow component (FWHM $\approx 80 \text{ km s}^{-1}$) and a broader component, larger than the clump circular velocity. When examined after a peak in the local SFR (caused by the swallowing of surrounding gas by the clump), a broad and irregular spectrum is obtained, corresponding to the more disturbed morphology of the same clump on the fourth panel of Figure 5.

We performed the same exercise for all clumps in our samples and display the median stacked spectrum on Figure 11, after rescaling the velocities by a factor $\sqrt{10^9/M_c}$ where M_c is the mass of the clump in solar masses, i.e., normalizing the baryonic mass of clumps to $10^9 M_{\odot}$ at all times. The characteristic stacked spectrum displays a narrow component and a broad component, well-fitted by the sum of two Gaussians of respective widths 93 and 424 km s^{-1} FWHM. This velocity widths are close to the circular velocity,¹⁷ and above the local escape velocity, respectively, for a normalized clump mass of $10^9 M_{\odot}$. The best fit for the stacked spectrum is obtained with 32% of the gas mass in the broad component.

We can confirm in our simulations that this broad component in the spectrum of clumps corresponds to a gas outflow, rather than infall. First, they are obtained with quasi-face-on orientation, so there is little fuel for infall onto clumps along the line of sight in our idealized experiments. Second, gas in the broad component has temperatures if the 10^4 – $5 \times 10^6 \text{ K}$ range, indicative of heating by stellar feedback processes. Third, the broad component *above* $\sim 200 \text{ km s}^{-1}$ typically contains $7 \times 10^7 M_{\odot}$ of gas mass, which is able to escape from a kpc-sized clump in $\sim 10^7 \text{ yr}$, consistent with our direct measurements of outflow rates at a few solar masses per year. This lends support to the idea that similar broad components in observed spectra of high-redshift giant clumps (e.g., Shapiro et al. 2009; Genzel et al. 2011; Newman et al. 2012a) would be most likely attributable to outflows. The velocity of the gas outflowing from the clumps scales with both clump mass and galaxy mass (Figure 12),

¹³ <http://youtu.be/Qm5-SkgnDYs>

¹⁴ Parallel to the initial disk plane.

¹⁵ Measurements across the spherical boundary may include diffuse gas that passes next to the clump and enters/leaves the boundary. Hence, the choice of the planar boundaries above and below the disk plan ensures to capture the feedback-induced outflow.

¹⁶ Equivalent to 0.07 arcsec FWHM at $z \approx 2$.

¹⁷ The FMHW of the narrow component is, in detail, somewhat lower than the circular velocity, which comes from the fact that these spectra were obtained for quasi-face-on galaxy orientations, and the giant clumps tend to be relatively aligned with the galaxy disk (Ceverino et al. 2012), which reduces the apparent face-on velocity amplitude.

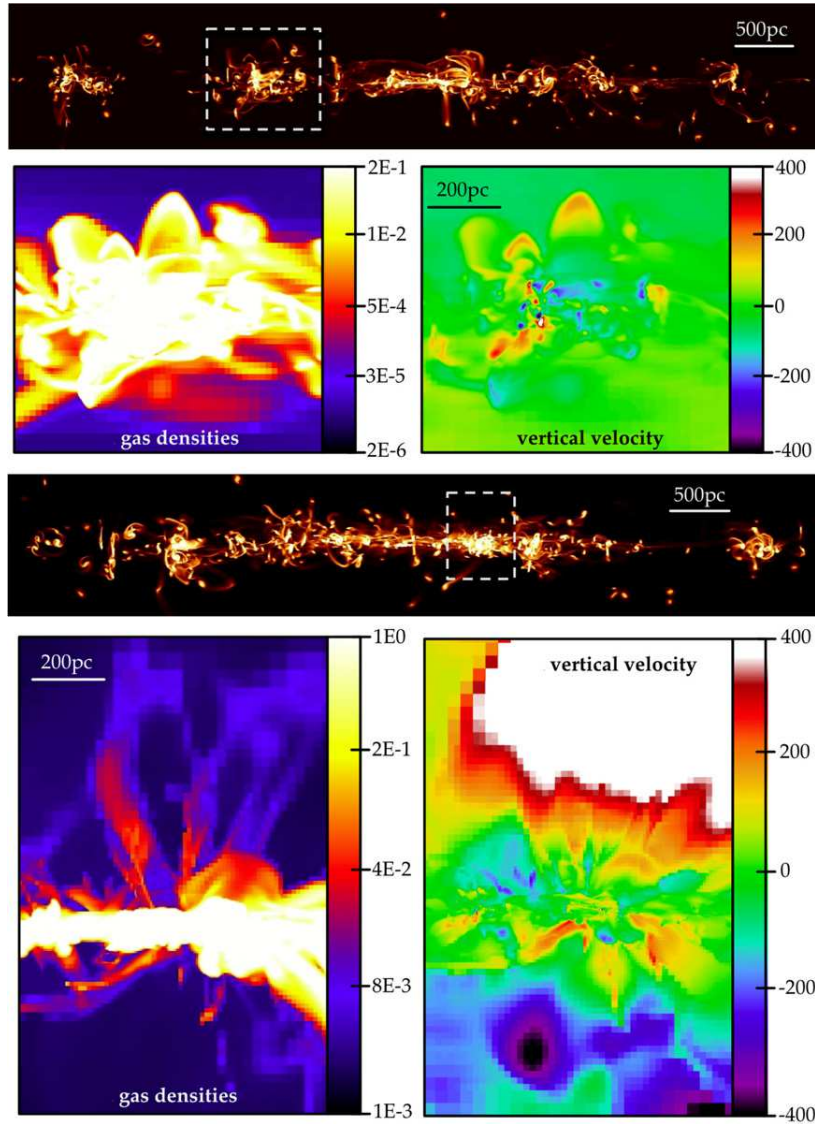


Figure 7. Birth of a gas outflow from a long-lived giant clump in model G'2. The entire galaxy is shown edge-on to the top (gas column density maps) after 90 and 270 Myr in the simulation, i.e., shortly after clump formation and at a more evolve stage. A giant clump is tracked and is shown in the inset maps of the mass-weighted average number density of gas (in cm^{-3}) and vertical velocity perpendicular to the disk plane (in km s^{-1}). At the early stage, shock fronts are formed by the outflowing gas onto the hot and diffuse halo surrounding the galaxy, with number densities of 10^{-2} – 10^{-1} cm^{-3} and velocities around 200 km s^{-1} . Later on a higher-velocity outflow escapes the clump region at velocities of 300 – 400 km s^{-1} with densities of a few 10^{-2} cm^{-3} , corresponding to an outflow rate of 1 – $2 M_{\odot} \text{ yr}^{-1}$ across a section of 1 kpc^2 . Note that the outflow continues to be accelerated at more than 1 kpc above the disk mid-plane, because of a pressure gradient, which turns the local outflow into a global galactic-scale outflow. A movie is available (<http://youtu.be/Qm5-SkgnDYs>) to show the development and expansion of gaseous outflows from clumps to galactic scales.

(A color version of this figure is available in the online journal.)

with a tighter relation when considering the entire galaxy mass. This shows that giant clumps can launch gaseous outflows that are rapid enough to escape the galactic potential well, even in high-mass galaxies. As previously said, more massive galaxies have more numerous and more massive clumps; and hence, the outflows initiated in the giant clumps can reach higher velocities there.

It is interesting to note that the observed velocity of gaseous outflows in Newman et al. (2012a), which are known to be largely launched by giant clumps (Genzel et al. 2011), scales with galaxy mass, which appears to be consistent with our results. More generally, the global properties of the outflows in our simulations (wind velocity, outflow rate compared with SFR, density range) are consistent with existing observational

constraints at $z \approx 0.5$ – 2 (Kornei et al. 2012; Rubin et al. 2013; Martin et al. 2013). These observations target high-redshift star-forming galaxies: these are not selected for clumpiness, but as a matter of fact these galaxies are generally very gas-rich, violently unstable, with significant clumpiness in 50%–70% of cases in the optical and near-infrared (Elmegreen et al. 2007; M. Mozena et al., in preparation; Y. Guo et al., in preparation). Our simulations are even consistent with the observed dependence of outflow velocities on galactic mass (Bordoloi et al. 2013). Hence, our feedback model produces realistic outflows, although the outflow velocity or mass loading are not imposed a priori in our feedback modeling, and indicates the star-forming clumps can be efficient launching sites for the outflows without being rapidly disrupted.

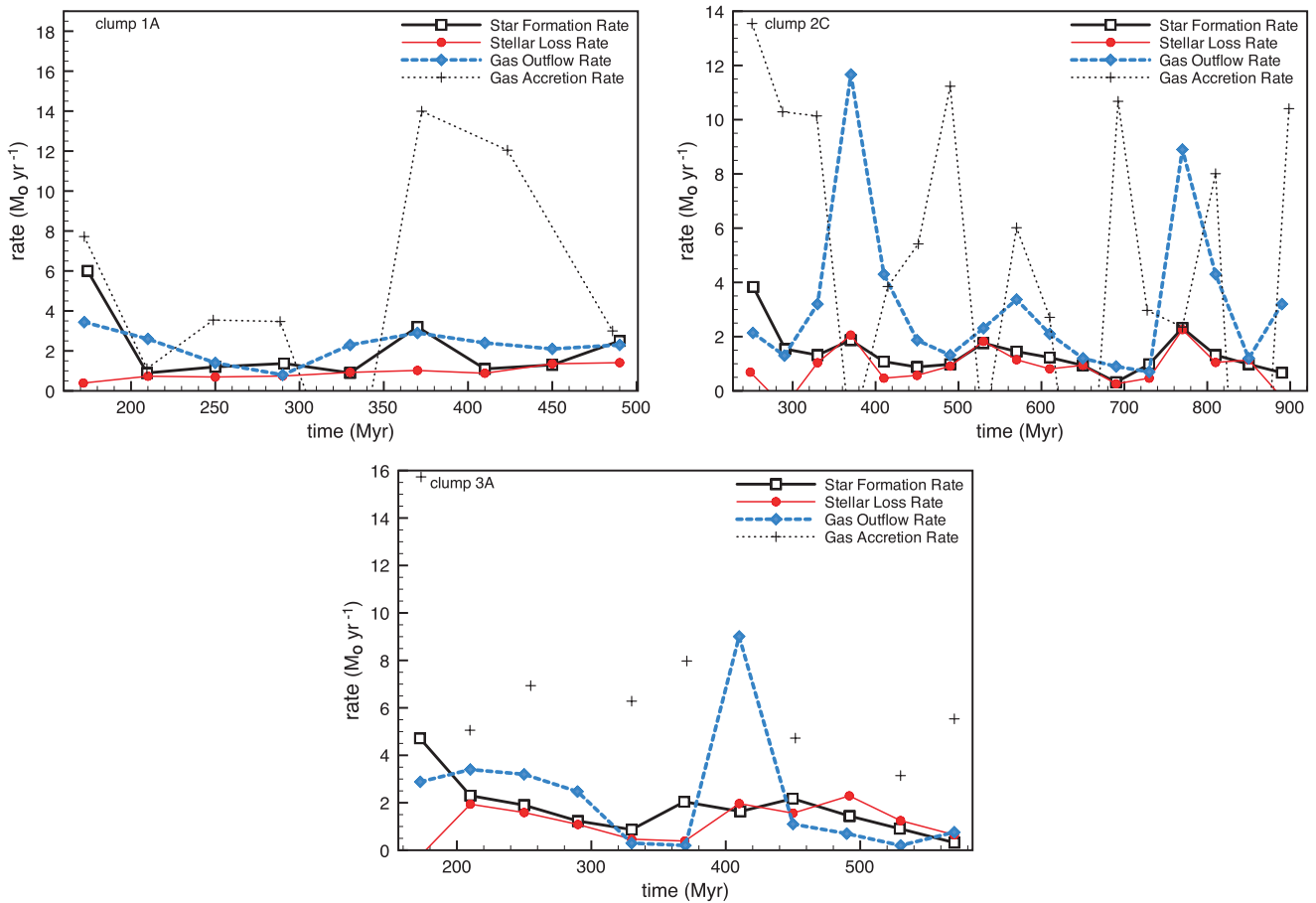


Figure 8. Time evolution of the SFR (thick solid line), outflow rate (thick dashed line), gas infall rate (thin dashed line), and stellar evaporation rate (thin solid line), for three representative clumps. We picked the longest-lived clump in each of the three fiducial simulations to increase the sampling of possible events. Gaussian smoothing of FWHM 40 Myr was applied to all quantities for clarity. The SFR and outflow rate typically increase after periods of intense accretion of the surrounding diffuse gas and/or smaller clumps. There is also an initial burst of star formation when the each clump initially collapses ($t = 0$ is the beginning of each simulation, and the first data points are 20 Myr after the first detection of each clump).

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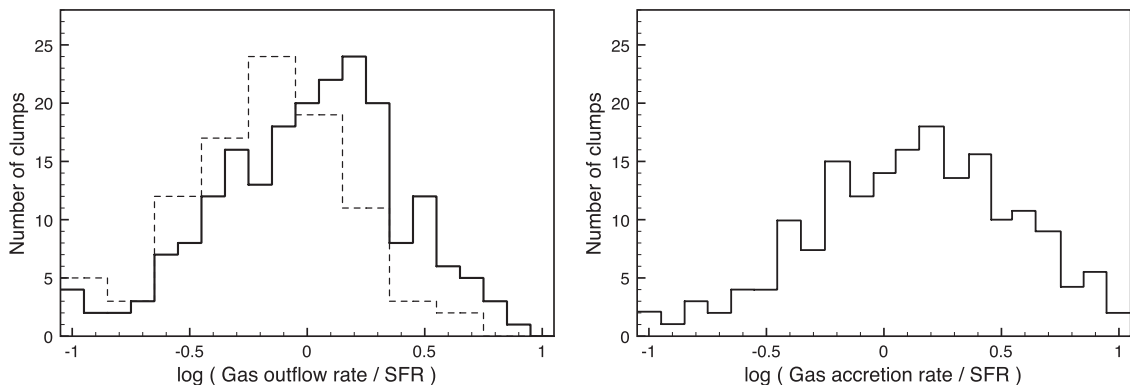


Figure 9. Left: statistical distribution of the outflow rate to SFR ratio at the scale of giant clumps. The thick histogram is for the cumulated distribution in the fiducial runs G1, G2, and G3. The dashed histogram is for run G'2 (normalized to the same maximum), in this run weaker supernovae feedback is employed along with the same radiation pressure model. The difference shows that the birth of outflows is not ensured solely by radiation pressure, but that supernovae explosions or at least their coupling with radiation pressure play a significant role. Right: rate of gas accretion by the clumps, for the fiducial runs G1, G2, and G3. It is of the order or slightly larger than the gas outflow rate.

3.3. Dynamical Loss of Aged Stars

The release of material by a clump is not limited to gas outflows. Tracking individual star particles throughout the simulations, we actually measure that a large fraction of stars captured by a clump during its initial collapse, or formed in situ

in the clumps, can gradually escape from the clump in a few 10^8 yr.

Independently from any feedback-driven gas outflows, it was already noted by Elmegreen et al. (2005) that typical clumps are not tightly bound compared with the tidal field of their host galaxy, with densities only a factor 10 above the limiting tidal

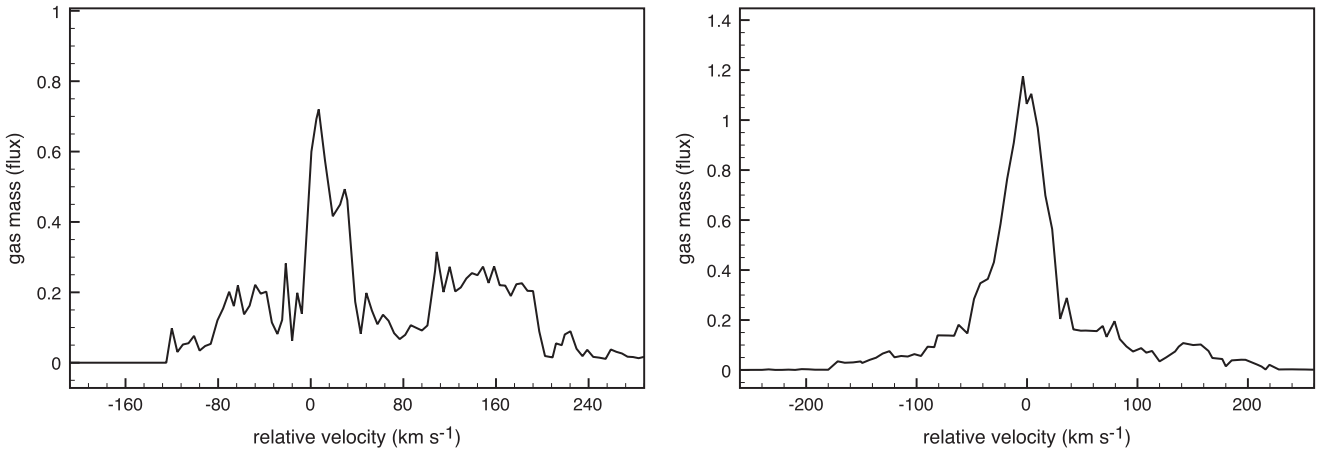


Figure 10. Line-of-sight velocity distribution (i.e., synthetic spectrum) of clump 2C, observed with a beam 600 pc FWHM, with an almost face-on orientation of the host galaxy. Left: the system is shown during its peak of outflow activity after a big gas cloud was absorbed (see Figure 5 and 8), the clump spectrum is broadly irregular with several high-velocity components. Right: we show the system in a calmer phase, 80 Myr later, when the mass outflow rate is marginally higher than the SFR. A double-Gaussian profile is observed, as in the stacked spectrum for all clumps shown in Figure 11.

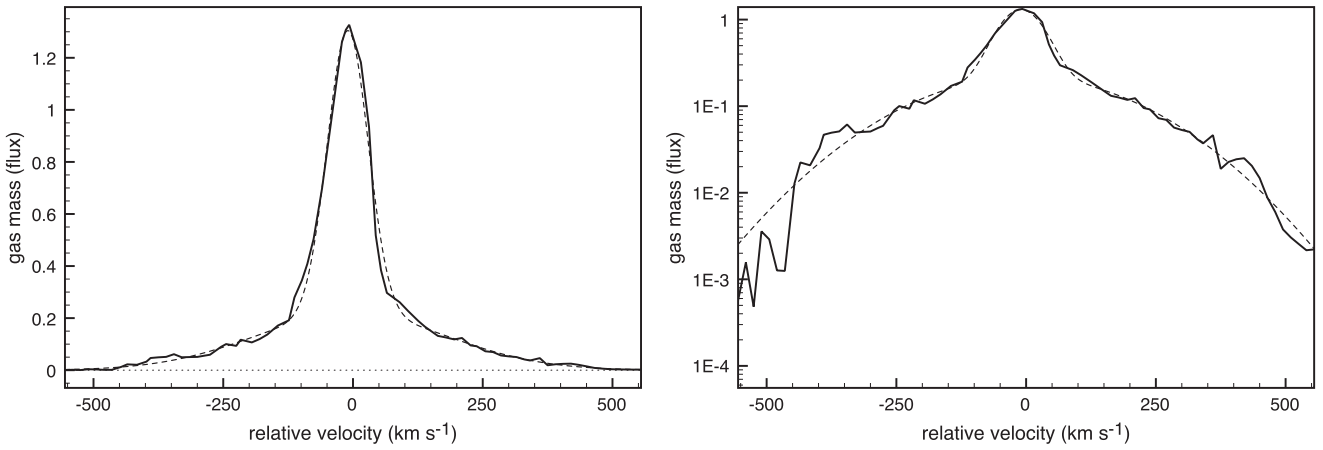


Figure 11. Median stacked spectrum comprising all the studied clumps (left: linear scale—right: log-scale), after rescaling each individual spectrum to the same clump mass (see text). The stacked spectrum is well fitted by a double Gaussian model (dashed), according to which the broad component contains 32% of the gas mass (a significant part of which, but not all, is above the clump escape velocity), 68% is in the narrow component (bound to the clump). In our simulations, the gas in the broad component is hot ($\geq 10^{4-5}$ K), outflowing gas.

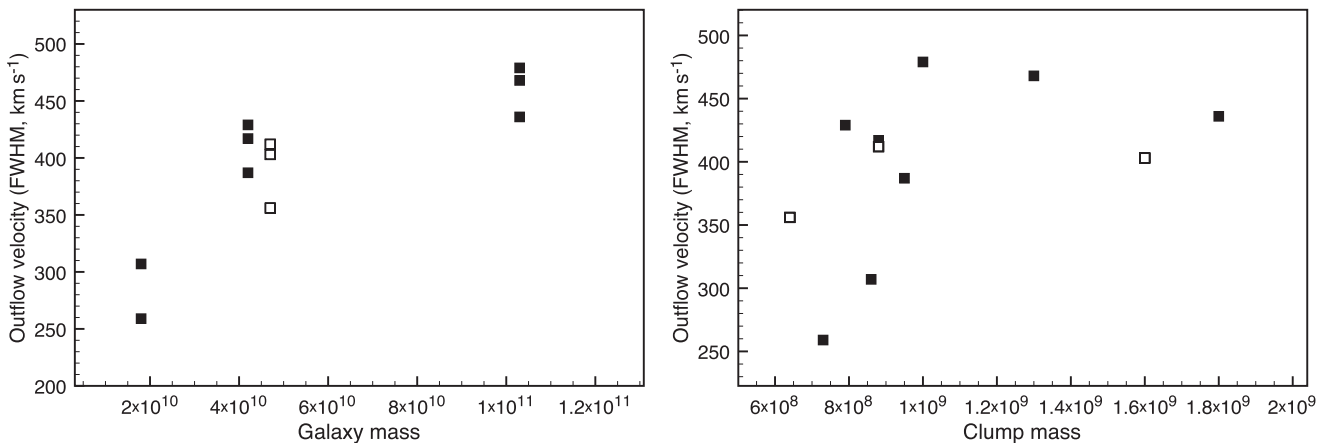


Figure 12. Outflow velocity for each clump, measured as the average FWHM of the broad spectral component for each individual clump (spectra are extracted every 40 Myr and stacked), as a function of galaxy mass (left) and average clump mass (right). The open symbols are for model G².

density, so that their outer parts could be significantly affected by tidal stripping. Previous simulations did find that stars formed in a clump can gradually escape the clump because of their increasing velocity dispersions over time and because of the galactic tidal field (see, for example, BEE07), and the released

stars were proposed to fuel the thick disk (Bournaud et al. 2009). The same effect is present in our simulations, in somewhat larger proportions, presumably because the gas outflows both regulate the growth of the clump mass and increase the stellar escape rate from the shallower local potential well.

To quantify the dynamical loss of stars by clumps, we measure the mass of stars present in the clump at a given time and that have left the clump later on.¹⁸ The fact that the obtained quantity is always positive confirms that it corresponds to the loss of in situ material and is not significantly affected by chaotic entry and escape of stars formed elsewhere. This typical “stellar loss rate” is $0.2\text{--}2 M_{\odot} \text{ yr}^{-1}$ for giant clumps, somewhat lower but of the order of their internal SFR. The clumps can release aged stars almost at the rate at which they form new stars.

To quantify the effect on the observable stellar ages of the clumps, we examine relatively old clumps in our simulations. For each of the eight clumps tracked in our fiducial runs, we “observe” it at the last snapshot before it enters the central kiloparsec, i.e., at a relatively advanced stage—the idea being that young clumps cannot have their stellar content severely affected by the long-term loss of stars. The “aged clumps” picked this way have ages ranging from 230 to 730 Myr, on average 400 Myr. Then, for each of these clumps, we identify the stars that have formed in the clump at any time, independently of their final location. We find that *the median age of stars that have formed in a given clump is about half of the real age of the clump* (Figure 13, dashed curves). This is because the clumps maintain a relatively constant star formation history owing to external gas accretion. Then, we identify the stars that are present in each clump at the “observed” time, and we find that because a higher fraction of aged stars have escaped from the clump, *the median age of stars lying in a clump is less than half of the real age of the clump* (Figure 13, solid curves). Our sample of “aged clumps” have a median real age of 420 Myr (ranging from 230 to 730 Myr), but the age of the stars that they contain has a median of 150 Myr (ranging from 120 to 210 Myr).

The effect is weaker for younger clumps. We show in Figure 14 the median age of the stellar populations contained in a clump as a function of the real clump age, as gathered from our entire sample. The median age of stellar populations saturates around 200 Myr even for older clumps, because the non-declining SFR is comparable to the gradual dynamical loss of older stars. Note that clumps do capture preexisting stars during their initial collapse, as expected for a two-fluid instability (Fellhauer et al. 2006; Elmegreen 2011); and thus, very young clumps can have a median stellar age older than the clump itself, as also seen in Figure 14.

We also learn that the stellar mass of clumps is regulated, as the rate of new star formation is not much larger than the stellar escape rate; thus, the stellar mass in a given clump does not strongly increase over time. The average stellar mass fraction in the clumps at 100–200 Myr of age is 18% (the remaining 82% is gas), and it only marginally increases to 23% in the 400–600 Myr age bin. Hence, the contribution of clumps to the stellar mass distribution is weaker than their contribution in the spatial distributions of gas and star formation. As illustrated in Figure 15, the contrast of clumps w.r.t. surrounding disk material in stellar mass density maps is low, typically less than a factor two. The clumps are much more contrasted in the equivalent optical light image (in which recently formed stars are more prominent), which compare favorably to real galaxies at high redshift (Figure 15).

¹⁸ We subtract from this quantity the mass of stars lying outside the clump at the former instant and found in the clump and the latter one, which means that the measurement is corrected for potential rapid entry/re-escape of preexisting background stars from the large-scale disk, which is found to be a relatively minor quantity.

3.4. Accretion Rate Onto the Clumps

Comparing the gas mass of clumps and the measured outflow rate, we infer the rate of gas accretion by the clumps, which would be hard to measure directly given the non-Lagrangian nature of the gas component in our simulations. The result is shown for several clumps in Figure 8. The obtained value is sometimes negative, which traces the uncertainty in other quantities and/or the potential tendency to slightly underestimate the outflows.¹⁹ The average of the positive values is five times larger than the absolute average of the negative values, confirming that the uncertainty is relatively small, and that we are measuring net inflow.

The gas accretion rate by the clumps is typically $2\text{--}15 M_{\odot} \text{ yr}^{-1}$, with the peak values corresponding to the accretion of smaller gas clouds (some of which are visible in the simulations snapshots), while the lower background values correspond to continuous accretion of diffuse disk material. This accretion rate is higher during the first 50 Myr, when each clump is initially collapsing (Figure 8), then it fluctuates about a level of a few $M_{\odot} \text{ yr}^{-1}$.

The accretion rates measured in our simulations are consistent with those predicted analytically by Dekel & Krumholz (2013). The predicted rate of capture of gas by a massive clump in a disk is $\alpha \rho_d R_{\text{tid}}^2 \sigma_d$ with $\alpha \approx 1/3$ (Dekel & Krumholz 2013), where ρ_d is the typical density of gas in the disk, σ_d the gas-velocity dispersion, and R_{tid} is the tidal radius beyond which gravitational forces from the galaxy dominate w.r.t. the internal gravity of the clump. We apply this to galaxy G2, where the measured gas velocity dispersion is $\sigma_d = 44 \text{ km s}^{-1}$, and the gas density in the mid-plane at the half-mass radius is $2.3 M_{\odot} \text{ pc}^{-3}$. The average mass of clumps 2A, 2B and 2C (averaged over the clumps and over time) is $7.8 \times 10^8 M_{\odot}$. The tidal radius at which forces from the galaxy (of mass²⁰ $1.2 \times 10^{11} M_{\odot}$) and from the clump are roughly equal²¹ is $R_{\text{tid}} \approx \sqrt{7.8 \times 10^8 / 1.2 \times 10^{11}} \times R_{\text{gal}}$, where $R_{\text{gal}} = 3.3 \text{ kpc}$ is the half-mass radius of this galaxy. This gives $R_{\text{tid}} \simeq 266 \text{ pc}$, so the predicted average accretion rate onto the clumps is about $7 M_{\odot} \text{ yr}^{-1}$. The average measured value in our simulation, after the first 100 Myr during which the clumps are collapsing with high infall rates, is $5.3 M_{\odot} \text{ yr}^{-1}$, very consistent with the simplified analytic estimate provided by Dekel & Krumholz (2013).

4. DISCUSSION

4.1. Observed Outflows and Stellar Populations

The recently detected outflows in giant clumps at $z \approx 2$ have gas outflow rates of $\approx 1\text{--}2 \times \text{SFR}$, with perhaps extreme cases up to $8 \times \text{SFR}$ (Genzel et al. 2011; Newman et al. 2012a). About 60% of the clumps in our models are in this range. Given that accurately resolving the clumps requires high sensitivity and adaptive optics, the observed samples remain very small and mostly consist of the brightest clumps in the most massive disk galaxies: their individual mass and SFR are typically higher than

¹⁹ As explained in Section 3.2, we chose to measure the outflowing component perpendicular to the disk, the outflow around spherical boundaries being somewhat larger, but consistent. This choice was made to be conservative in the estimated outflow rates. The negative values of the estimated gas accretion rate onto the clumps can thus correspond to gas that leaves the clumps within the disk plane, either because of in-plane feedback-driven outflows, or because of tidal stripping of the gaseous component.

²⁰ Including the mass of dark matter inside the galaxy half-mass radius.

²¹ If we approximate both the clump and the galaxy with two point masses on a circular orbit.

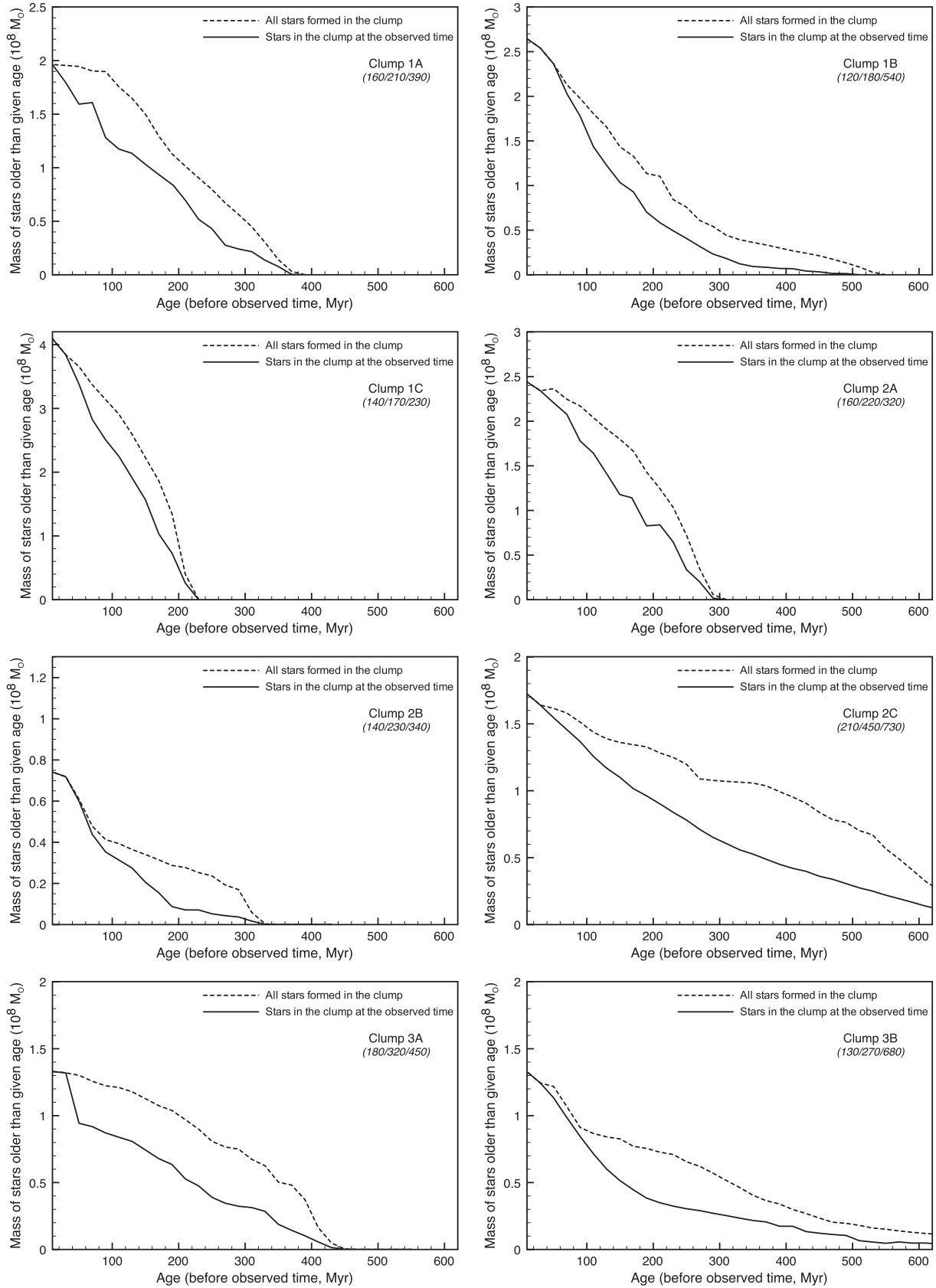


Figure 13. Cumulative age distribution of stars present in clumps, for a selection of relatively old clumps (see text). The dashed line shows for comparison the cumulative age distribution of all the stars that formed in each given clump, regardless of their location (in or outside the clump) at the analyzed instant, and normalized to the same final value for clarity. Three numbers are given in Myr for each case: median age of the stars that lie in the clump at this instant/median age of the stars that have formed in the clump (but may not lie in the clump anymore)/actual age of the clump main progenitor as tracked in the simulation. The typical age of stars in an evolved clump is rarely larger than 200 Myr, even for clump ages of 300–600 Myr and more.



Figure 14. Median age of the stars contained in the clump, as a function of the clump age. We show the average measurement, for all clumps in each age bin of 100 Myr. Clumps of 500 Myr and more contain stars of typical age ≤ 200 Myr, because of the dynamical loss of older stellar populations. Young clumps of 100 Myr and less can have an average stellar age larger than the actual clump age, because they capture preexisting stars during their initial collapse. We assumed random ages from 0 to 1 Gyr for the stars present at the beginning of the simulation.

in the more “typical” clump selected in our models, which might lead to different trends in the outflow rate to SFR ratio (stronger activity, but deeper potential wells). Given these limitations, and despite the strong agreement between our fiducial models

and observations, it is probably premature to assume that our alternative model with reduced supernova feedback, which typically produces outflow rates somewhat lower than the SFR, is less realistic. Nevertheless, the comparison of this model with our fiducial ones shows that supernovae do play a role in clump-scale outflows, the rate of which is not determined solely by the momentum injection through other feedback modes.

Our simulations are also in agreement with the observed stellar populations. The continuous accretion of mass by the clumps keeps their gas mass about constant, and their star formation histories are relatively flat (with large fluctuations, but without a significant systematic trend). Wuyts et al. (2012) indicate a typical age for stellar populations of 100–200 Myr, assuming exponentially decaying star formation histories. This assumption can somewhat underestimate the age of stellar populations with more constant SFR histories (Maraston et al. 2010). The continuous re-formation of young stellar populations in the clumps, coupled with the gradual tidal loss of aged stars, make the clumps of real age 300 Myr and above appear as containing a median stellar age of about 200 Myr. This effect, already noted²² in Bournaud et al. (2007), puts our simulations in agreement with the typical observed stellar ages in the clumps.

²² The effect was somewhat weaker in the Bournaud et al. (2007) models, presumably because the lack of gaseous outflows maintained a deeper potential well that could retain the stars on longer timescales.

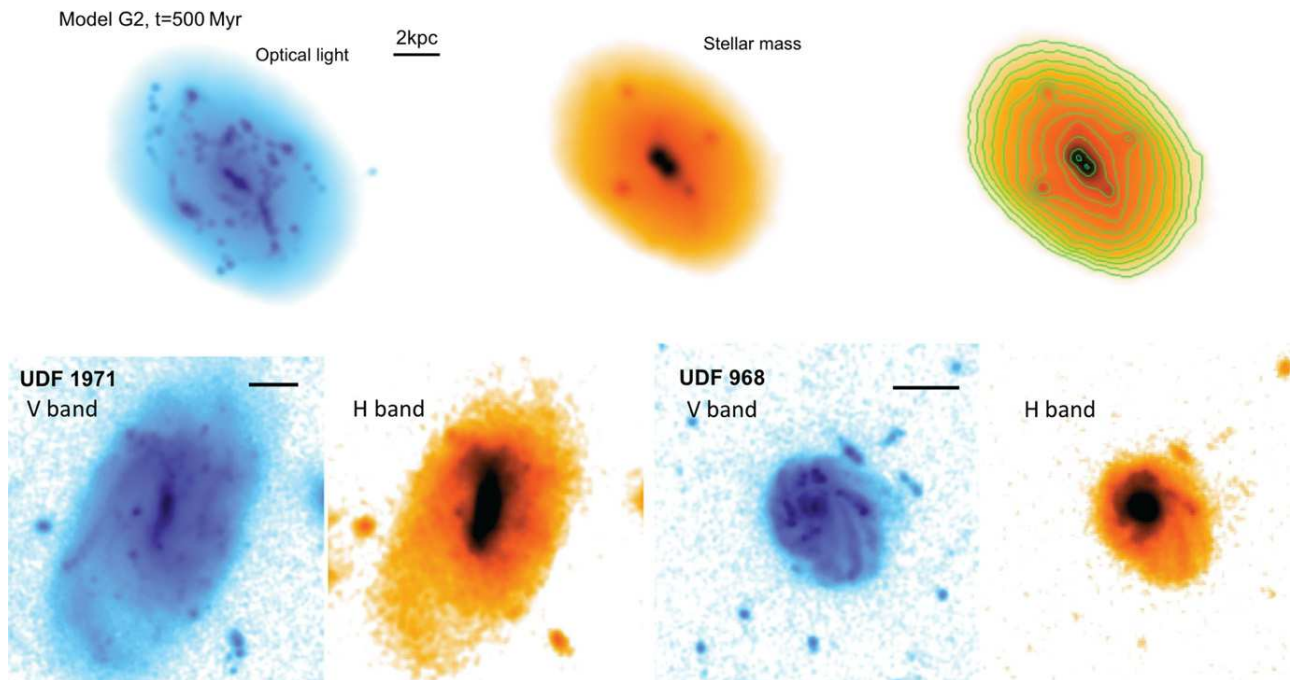


Figure 15. Stellar mass map and optical light image of galaxy G2 at $t = 500$ Myr. The optical light image is generated assuming that the mass-to-luminosity ratio of stellar populations is constant during 10 Myr, and subsequently decreases as $t^{-0.7}$. The contrast clump in stellar mass maps is much lower than in optical images or gas density maps. For example, the clump visible to the top is clump 2C, which has formed 400 Myr ago in the outer disk. On the right panel, logarithmic isodensity contours with a spacing factor of 1.5 between contours are overlaid (i.e., the level of a given contour is 1.5 times the level of the previous one). It shows that on such stellar mass maps, the peak surface density of a clump is only 1.5–2 times that of the surrounding material. Here, we applied Gaussian smoothing with FWHM of 200 pc, lower than the clump size, and no noise was added, so the low contrast of the clumps is not an effect of dilution at low resolution. Note that the typical resolution of *Hubble Space Telescope* (HST)/WFC3 imaging may lead to some extra decrease in the clump contrast because of lower-resolution beam smearing. On the bottom, we show the *HST*/ACS and *HST*/WFC3 observations of two clumpy galaxies from the Elmegreen et al. (2007) sample, UDF 1971 and UDF 968. The clumps have a low contrast in the near-infrared but the optical images are much more clumpy with giant clumps, smaller clouds, some spiral features, and offset centers, similar to the advanced stages in our clumpy galaxy simulations. Note that the small knots seen around these two UDF galaxies are not minor mergers or satellites, but background galaxies (at least for most of them, the photometric redshift is much larger). The typical range of clump masses measured by Elmegreen et al. (2007) is consistent with our current set of models.

(A color version of this figure is available in the online journal.)

4.2. Short-lived Clump Models

Other numerical simulations of high-redshift clumpy disks have used strong momentum-driven feedback to model short-lived clumps (Genel et al. 2012b; Hopkins et al. 2012). In addition to using a high trapping factor for photons, providing high momentum to the gas, these models assumed that the initial outflow velocity should be equal to the local escape velocity from the galaxy or from the clump (multiplied by a scaling factor close to unity). This hypothesis maximizes the deduced mass loading of the outflows. For a given momentum $m \times v$, this puts the highest possible gas mass m above the escape velocity. The observed wind velocities may actually be larger (Genzel et al. 2011), which would reduce the involved mass for a given amount of momentum. In addition, the hydrodynamic interaction was also suppressed over some distance in the Genel et al. (2012b) models, preventing dilution of the outflow momentum and thus boosting up the mass loading. In contrast, in our simulations, we evaluate the outflow properties of velocity and mass loading from the physical elements of the feedback processes, and we find that some components in the outflows are much more rapid than the escape velocity (up to 500 km s^{-1} in agreement with data), in which case the resulting mass loading is lower for the available momentum. In contrast, some regions in the outflows have moderate velocities (around 100 km s^{-1}) and remain eventually bound in the clump’s potential well unless they exchange momentum with other components.

These other simulations with very extremely strong feedback have demonstrated that it is possible to find combination of star formation efficiencies and feedback schemes that allow clumps to gather about 10^9 solar masses of baryons for a short time before clump disruption. However, the arbitrarily strong feedback imposed disrupts the clump within typically 30–50 Myr, which appears very short compared with observed stellar populations ages up to 200 Myr, and other clumps do not re-form on an equally short timescale. Hence, it is rather unlikely that short-lived clump models could account for the high frequency of giant clumps in high-redshift disks and the 200 Myr-old stellar populations found in these clumps.

4.3. Long-lived Clump Models

The direct interpretation of the models presented here is that the inclusion of radiative feedback, and the observations of outflows launched by giant clumps and relatively young stellar ages in the clumps, are fully consistent with long-lived clumps that keep their mass roughly constant about the initial Jeans mass, drive an inflow throughout the disk, and migrate radially, promoting bulge growth (in realistic proportions, as further discussed in the next subsection). It does not mean that this long-lived clump model is the most realistic one: models with stronger feedback and short-lived clumps remain a potentially viable alternative.

A full parametric study of feedback and the detailed dependence of clump lifetime on feedback efficiency is beyond the scope of the paper, given the cost of simulations with the spatial and mass resolution used here. We nevertheless note that we have varied the timescale for the dissipation of energy injected by supernovae, and the resolution, without reaching significant variation in the clump properties and evolution (comparing in particular models G1 and G’2). We also note that other studies using very different feedback schemes also find realistic outflows with long-lived clumps (e.g., Perez et al. 2013), suggesting that long-lived clumps are the natural outcome over a broad

range of feedback parameter. At the opposite short-lived clumps have been obtained so far only with very strong radiative feedback attributed to very efficient photon trapping in the dense gas clouds. Such high trapping factors are not favored by systematic study of radiative feedback in cosmological simulations, because they lead to the quasi-absence of thin disks in present-day spirals, and also lead to strongly deficient amounts of star formation below $z \approx 2$, especially compared with those achieved at higher redshift (Roškar et al. 2013). At the opposite, moderate photon trapping factors of about a few units, as assumed in this study, are found to result in star formation histories consistent with the plausible range supported by observations, and to remain consistent with the presence of thin star-forming disks in $z = 0$ spirals. This suggests that the feedback parameters used in this study are more realistic, yielding further (but indirect) support to the long-lived clump picture. Note also that the ejection of gas from the clumps and the resulting loss of aged stars occur at realistic rate in our current simulations, resulting in a range of stellar ages in the clumps very consistent with the observations. Simulations with increased star formation and feedback efficiencies will be presented in V. Perret et al. (in preparation) showing that stronger feedback increases the rates of gas ejection and reaccretion and of associated stellar loss, yielding stellar ages in clumps systematically younger than the observed range. This is another indication that the current model could be the most realistic solution, given the existing theoretical and observational constraints.

4.4. The Regulated Growth of Bulges

The inclusion of realistic outflows in our current model induces several interesting differences compared with previous long-lived clump models. The baryonic mass of clumps is, on average, constant over their inward migration in the disk. Models without feedback or limited to weak supernova feedback found clumps whose mass was steadily increasing along their inward migration, because of the same process of continuous mass accretion as in the current models (e.g., Noguchi 1999; Bournaud et al. 2007, see also the clump growth in the “no wind” models by Genel et al. 2012b). The continuous mass accretion by clumps can be largely compensated by gaseous outflows²³ in our new models, and as a result the mass of clumps when they reach the galactic center is regulated to 0.2%–2.4% (on average 0.8%) of the galaxy baryonic mass in our current simulation sample. Furthermore the loss of aged stars and reaccretion of gas keeps much of this mass gaseous, with an average gas fraction in the clumps in our sample of about two thirds, and an average stellar mass of $3 \times 10^8 M_{\odot}$.

If a generation of clumps contains 5–10 clumps, each conveying $3 \times 10^8 M_{\odot}$ of stars to the bulge in ≈ 500 Myr, and three such generations span a galaxy’s life in the $z \approx 3$ –1 range, this mechanism overall provides about $7 \times 10^9 M_{\odot}$ of stars to a bulge component—which would be a “classical” bulge for a clump coalescence mechanism (Elmegreen et al. 2008a). This is a rather reasonable mass for galaxies with stellar masses reaching several $10^{10} M_{\odot}$ by redshift zero. The clumps will also carry gas inward (through their migration if they are long-lived, and through the general instability-driven inflow in any case) but if the gas preserves a significant angular momentum it may form the inner regions of exponential disks, or a disky (pseudo-)bulge (see also Inoue & Saitoh 2012). Hence, there is not necessarily

²³ Stellar loss was present in previous models, but probably weakened by the deeper potential wells associated to the absence of gaseous outflows.

over-production of bulges in such a long-lived clump model. Alternatively, the instability-driven inflow may grow a spheroid in which central star formation is continuously fuelled with the appearance of “blue nuggets” (A. Zolotov et al., in preparation; Barro et al. 2013). A detailed budget of bulge growth and disk survival in our current clumpy disk models, and in similar models of high-redshift mergers, will be presented elsewhere (V. Perret et al. 2013, in preparation).

Many of the observed high-redshift clumpy galaxies are already more massive at $z \approx 2$ than today’s Milky Way, and are not expected to remain disk-dominated until redshift zero. Smaller galaxies, with the typical mass expected for a Milky Way like progenitor at $z \approx 2$, also seem clumpy in their morphology (even perhaps down to lower redshifts Elmegreen et al. 2007; Bournaud et al. 2012), and their apparently “dispersion-dominated” kinematics are also typical for clumpy disks once studied at high resolution (Newman et al. 2012). It appears in our simulations that such lower-mass galaxies tend to have somewhat lower-mass clumps. Over a 1 Gyr long period, our low-mass model G3 forms only four clumps more massive²⁴ than $4 \times 10^8 M_\odot$, while 12 such clumps are found in the high-mass model G1. Given that lower-mass clouds are easily disrupted by feedback (and that the associated outflows may more easily escape the galactic potential well), the relative contribution to bulge growth should be weaker in low-mass galaxies. This could naturally explain the correlation between stellar mass and bulge fraction along the Hubble Sequence, but should be studied with physically motivated feedback models in full cosmological context so as to encompass the full duration of clumpy phases. It is observed that lower-mass galaxies can form clumps at lower redshifts, but with more modest clump masses (Elmegreen et al. 2013).

The bulges that grow in our models, in addition to being reasonably massive, are very gassy, because the clumps evolving with realistic feedback remain gas-dominated—simulations with weak supernovae-only feedback produce clumps whose mass and stellar fraction were increasing over time. Note that Perez et al. (2013) also studied the evolution $z \approx 2$ unstable disks, using on a fairly different stellar feedback model, also producing realistic outflows and wind mass loading along with long-lived clumps, and also conclude that the growth of bulges occurs in reasonable proportions considering the stellar mass of the involved galaxies. Although we defer the analysis of the structural properties of bulges in our current model to a subsequent paper, we here speculate that these bulges could often resemble pseudo-bulges, since they form as gas-rich central mass concentrations. Relatively stochastic growth of a pseudo-bulge could thus likely take place in these systems. The instability-driven inflow can grow the central supermassive black hole (SMBH) in a more continuous way, as predicted in simulations and supported by the high AGN fraction in clump disks (Bournaud et al. 2012) and the high AGN fraction in high-redshift star-forming galaxies in general (Mullaney et al. 2012; Juneau et al. 2013). Steady SMBH feeding along with more stochastic (pseudo-)bulge growth could thus be a path to explain the poor correlation between SMBH mass and pseudo-bulges (Kormendy et al. 2011). Note that if an intermediate-mass or SMBH lies inside a clump, it may undergo rapid²⁵ mass segregation and lie closer to the clump center than the stars, on

average, and therefore not be as likely to evaporate or leave the clump as the stars do. Such black holes could potentially migrate to the central bulge and SMBH (Elmegreen et al. 2008b), promoting a high SMBH to stellar mass ratio there if bulge growth is more regulated by escape of aging stars from the migrating clumps.

4.5. Self-regulation of the Clump Mass

We have noted in the previous parts that when the SFR of a given giant clump increases, its outflow rate also increases, and generally in larger proportions (at least when averaged over time periods of 40 Myr, Figure 7). We here suggest that this response can compensate for episodes in which the mass of a giant clump tends to grow, for example, when a smaller clump is swallowed. The strong subsequent outflows can prevent growth of the clump mass. Without this, the ability of clumps to accrete surrounding material would increase and their mass would continue to increase.

For an SFR = S expressed in $M_\odot \text{ yr}^{-1}$, the outflow rate can be written as S^q and the best fitting exponent in our models is found²⁶ to be $q \approx 2.2$. This is a crude description of the feedback behavior, but it captures the fact that moderate increases in the SFR can turn into larger increases in the outflow rate, as seen in the individual tracks on Figure 8. In a steady state the accretion rate should be $A \approx (1 + \epsilon_{\text{escape}})S + S^q$ to keep the clump mass constant, where $\epsilon_{\text{escape}} \approx 0.5$ is the fraction of stars formed in the clump that eventually escape the clump (after a time delay of the order of 200 Myr in our simulations, neglected here). A clump forming in a non-turbulent medium would have a mass accretion rate equal to the Jeans mass divided by the gravitational freefall rate, during its initial collapse phase. This quantity is $A = \sigma^3/G$. Since the giant clumps of high-redshift galaxies move rapidly in a turbulent disk, the surrounding reservoir is continuously replenished at the ambient disk density, keeping this accretion rate constant. The steady state condition is thus $(1 + \epsilon_{\text{escape}})S + S^q = \sigma^3/G$ with $q = 2.2$. Hence, a typical clump forming stars at a rate of $2 M_\odot \text{ yr}^{-1}$ can live in steady state if its binding velocity dispersion is 38 km s^{-1} , which is close to the measured dispersions in our models (Section 3.1), and in the molecular gas of high-redshift galaxies. The somewhat higher dispersions measured in the simulations ($38\text{--}53 \text{ km s}^{-1}$) may result from the tidal stirring of clumps, otherwise the excessive binding would lead to a steadily increasing clumps mass. Modeling the feedback response with a power law as done here is of course arbitrary, but the global idea is that there is a steep response where a moderate increase in the accretion rate and SFR can be followed by a stronger increase in the outflow rate (as previously see in Figure 8), leading to self-regulation of the clump mass.

5. SUMMARY AND IMPLICATIONS

We have presented simulations of high-redshift clumpy disk galaxies with a new implementation of photoionization and radiation pressure feedback, in addition to supernova feedback, and a physically motivated estimation of the wind mass loading. Previous models had considered only weak supernovae feedback models or had used strong feedback recipes generating high-velocity outflows with ad hoc high-mass loading factors. This paper has focused on the properties and evolution of the giant clumps of $10^8\text{--}9 M_\odot$ typically found in these galaxies, and the associated birth of outflows. The main findings are as follows:

²⁴ Staying above $4 \times 10^8 M_\odot$ for at least 40 Myr.

²⁵ The mass segregation timescale would typically be about 1–10 Myr for a $10^{5\text{--}6} M_\odot$ black hole in a giant clump.

²⁶ Best fit for the SFR and gas outflow rate averaged over 40 Myr time bins.

1. The physically motivated model for photoionization and radiation pressure feedback from Renaud et al. (2013), used either with a simple thermal model for supernovae feedback or the non-thermal model from Teyssier et al. (2013), produces massive gas outflows from the giant clumps, with outflow rates at the scale of individual clumps of the order of the SFR and up to 10 times the SFR over the short periods, where the SFR in a clump is $1\text{--}2 M_{\odot} \text{ yr}^{-1}$. The outflow properties are consistent with observations. Comparison of various models indicate that the outflow rate is not determined solely by radiation pressure, but is also influenced by the energy injection from supernovae.
2. The giant clumps survive for several 10^8 yr, until they migrate through torques and dynamical friction and coalesce centrally after a few hundred Myr. Their mass remains about constant in spite of the gaseous outflows. This is because the clumps are wandering in a gas-rich turbulent disk, from which they constantly accrete gas at a rate of a few $M_{\odot} \text{ yr}^{-1}$, compensating for the mass loss. Feedback-driven outflows ensure regulation of the clump mass after periods of high accretion rates.
3. The continuous accretion of gas sustains long-lasting star formation histories with relatively constant SFRs. The average age of stars in a giant clump is significantly younger than the age of the clump itself. In addition, the stars that form in a clump gradually leave the clump through tidal stripping, and up to half of the stars formed in situ at a given time can have left the clump 200 Myr later. This preserves young stellar populations (≤ 200 Myr), even in older clumps.
4. Bulge growth through instability-driven inflow, clump migration and coalescence, occurs in reasonable proportions on the current long-lived clump models, especially because the mass of clumps is self-regulated. Clump migration brings stellar mass to the central bulge at an average rate of 1–10 solar masses per year, over a timescale of 500 Myr for a single generation of clumps or more if cosmological accretion helps to re-form clumps. This typically grows a bulge of $10^9\text{--}10 M_{\odot}$ for present-day galaxy masses of $10^{10\text{--}11} M_{\odot}$, which is a reasonable bulge fraction. Furthermore, given the high fraction of gas in the material conveyed by the clumps, a large fraction of the central mass may evolve into a disky pseudo-bulge. As for more massive clumpy galaxies at $z \sim 2$, many of them evolve into ellipticals by $z = 0$ and the instability-driven and clump-driven spheroid growth may be an important ingredient in this evolution.

The observed clump properties in $z = 2$ galaxies (outflows and stellar content) are fully consistent with this long-lived clump model. Hence, the clumps, even if they represent a limited fraction of the total stellar mass, and more generally the associated violent disk instability, can have a major dynamical influence as studied in previous works: migration to the central bulge, fueling of bright AGN phases up to the Eddington rate (Gabor & Bournaud 2013, inside-out growth of a thick exponential stellar disk (Bournaud et al. 2007), and perhaps also globular cluster formation (Shapiro et al. 2010) and dark matter cusp erosion (Inoue & Saitoh 2011). Our feedback model shows that observations of outflowing gas and young stellar populations do absolutely not imply that clumps have to be short-lived transient features. However, since the strength of stellar feedback, and in particular the efficiency of photon trapping in the ISM, remains debated, it remains interesting to consider

models with extreme strong feedback (Genel et al. 2012b; Hopkins et al. 2012) as a potentially viable alternative.²⁷ These models differ by two aspects. First, they assume a very efficient trapping of the photons emitted by young stars, maximizing the feedback efficiency beyond theoretical expectations (Krumholz & Thompson 2012; Dekel & Krumholz 2013). Second, they impose the wind velocity to be equal or close to the local escape velocity, which maximizes the mass loading of the winds. Using a different approach to estimate the mass loading of momentum-driven winds from physical considerations (Renaud et al. 2013) we found here that the wind velocity can be higher, reducing the involved gas mass for a given amount of available momentum. Observations of a radial gradient in the stellar age of clumps (Förster Schreiber et al. 2011; Guo et al. 2012) suggests that real clumps do live long enough to migrate radially. Given that the gradient in real clump age should be steeper than in stellar age (Figure 14), these data do actually suggest lifetimes of a few 10^8 yr. Further support to the long-lived clump picture is provided by the parametric study of radiative feedback by Roškar et al. (2013), showing that very high trapping factors of photons in dusty clouds, which seem required to make giant clumps short-lived, are inconsistent with the observed star formation histories and the survival of thin gaseous disks in nearby spirals (see discussion in Section 4.3). At the opposite, long-lived clumps are obtained with a variety of feedback schemes producing realistic outflows (see, for example, Perez et al. 2013 and this study).

The fact that clumps continuously lose mass and reaccrete gas does not mean that they are permanently disrupted and reformed—as illustrated in our time series, the clumps are persistent structures that are permanently present with a roughly constant mass. When a clump is analyzed at a given time and 100 Myr later, it still contains about 80% of its initial baryons. Over a lifetime of 300–500 Myr, a typical clump will lose about 50%–65% of its initial baryonic mass (through gas outflow and dynamical loss of stars), and will reaccrete a comparable amount of gas. The clumps in unstable disks have a wave-like behavior like spiral arms in lower-redshift galaxies, with the interesting analogy that spiral arms have a very weak contrast in the near-infrared and in stellar mass maps, and mostly contain very young stellar populations and unvirialized molecular clouds, but are density waves that have propagated in the disk for timescales of 10^9 yr or more, with important dynamical impact (e.g., Puerari et al. 2000). A noticeable difference is that the timescale for mass exchange is longer for the high-redshift giant clumps, and complete renewal of their mass content does not occur faster than clump migration to the galaxy center.

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²⁷ Eventually, unambiguous determination of the effects of photon trapping would require full radiative transfer hydrodynamic simulations with sufficient resolution to resolve the size of young ionized regions in dense gas, i.e., parsec-scale resolution.

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4.4 Coalescence des complexes granulaires

La coalescence des complexes granulaires dans les simulations de l'échantillon MIRAGE s'est révélée être un mécanisme surprenant à bien des égards. Les granules atteignant le centre de masse où se trouve un progéniteur de bulbe stellaire subissent un champ de force de marée extrêmement élevé dans les derniers parsecs avant la fusion totale des granules. De manière similaire au mécanisme de fusion de galaxies, les noyaux des deux granules fusionnent tandis que les parties externes se trouvent accélérées par la force de marée. Une partie de la matière accélérée est en mesure de quitter le disque. Ce qui rend ce processus original dans les simulations de l'échantillon MIRAGE, c'est sa capacité à éjecter des volumes considérables de gaz en dehors du rayon optique de la galaxie hôte, avec des vitesses permettant d'atteindre le rayon de Viriel. Ce processus est illustré dans la figure 4.22 pour la simulation G1_G1_90_90_180. On constate que ce mécanisme éjecte de grandes quantités de gaz et modifie la structure du bulbe. Le processus de fragmentation d'un disque stellaire et gazeux, et la migration des granules vers le bulbe est un phénomène modélisé depuis plusieurs années (Bournaud et al., 2007a; Elmegreen et al., 2008; Inoue et Saitoh, 2011), mais une telle éjection de gaz massive associée à la coalescence n'est rapportée nulle part dans la littérature. Dans les parties suivantes, je décris quantitativement les effets de la fusion de granules, et je tente d'apporter les éléments permettant d'expliquer l'origine de ce processus. Tout d'abord, j'introduis un jeu de simulations complémentaires qui me permet de comparer avec le cas où les granules les plus massifs sont détruits par la rétro-action stellaire. Puis, j'introduis l'algorithme de détection des granules dans les simulations, et je quantifie statistiquement leurs propriétés. Enfin, j'étudie l'effet de la coalescence des granules de manière distincte sur le gaz, les étoiles, et le halo de matière sombre.

4.4.1 Simulations complémentaires

Afin de comprendre le rôle des complexes granulaires dans l'évolution de nos modèles de galaxies, j'ai réalisé des simulations complémentaires à celles de l'échantillon MIRAGE. Ces simulations, dont les conditions initiales sont strictement identiques, ont un traitement de la formation stellaire différent. La loi de formation stellaire s'écrit :

$$\dot{\rho}_\star = \begin{cases} \epsilon_\star \rho_{gas} / t_{ff} & \text{si } \rho_{gas} \geq \rho_0 = 100 \text{ cm}^{-3} \text{ et } T - T_{poly}(\rho_{gas}) < \Delta T_{poly} = 2 \times 10^5 \text{ K} \\ 0 & \text{sinon} \end{cases} \quad (4.2)$$

avec $\dot{\rho}_\star$ le taux de formation stellaire local, $\epsilon_\star=0.01$ l'efficacité de formation stellaire, ρ_{gas} la densité du gaz, $t_{ff} = \sqrt{3\pi/(32G\rho_{gas})}$ le temps caractéristique de chute libre dans la cellule AMR, ρ_0 le seuil de densité de formation stellaire, T la température de la cellule, $T_{poly}(\rho_{gas})$ la température du polytrope calculée pour la densité de gaz de la cellule AMR et ΔT_{poly} l'écart de température maximum avec le polytrope (cf. figure 1 de la section 3.2). Le polytrope de Jeans est introduit afin de résoudre au moins quatre fois la longueur de Jeans pour chaque cellule AMR et prévient ainsi la fragmentation numérique. La seule différence entre les simulations P et les simulations MIRAGE est que, dans ces dernières, la formation stellaire n'est pas autorisée au-dessus de la limite absolue $T = 2 \times 10^5$ K (cf. équation 1 de la section 3.2). Cette limite a pour effet de bloquer la formation stellaire dans les régions où le gaz se trouve sur la partie haute du polytope de Jeans : à partir de $\rho_{gas} \sim 3500 \text{ cm}^{-3}$, le polytrope de Jeans chauffe artificiellement le gaz à des températures supérieures à 2×10^5 K, ce qui signifie que

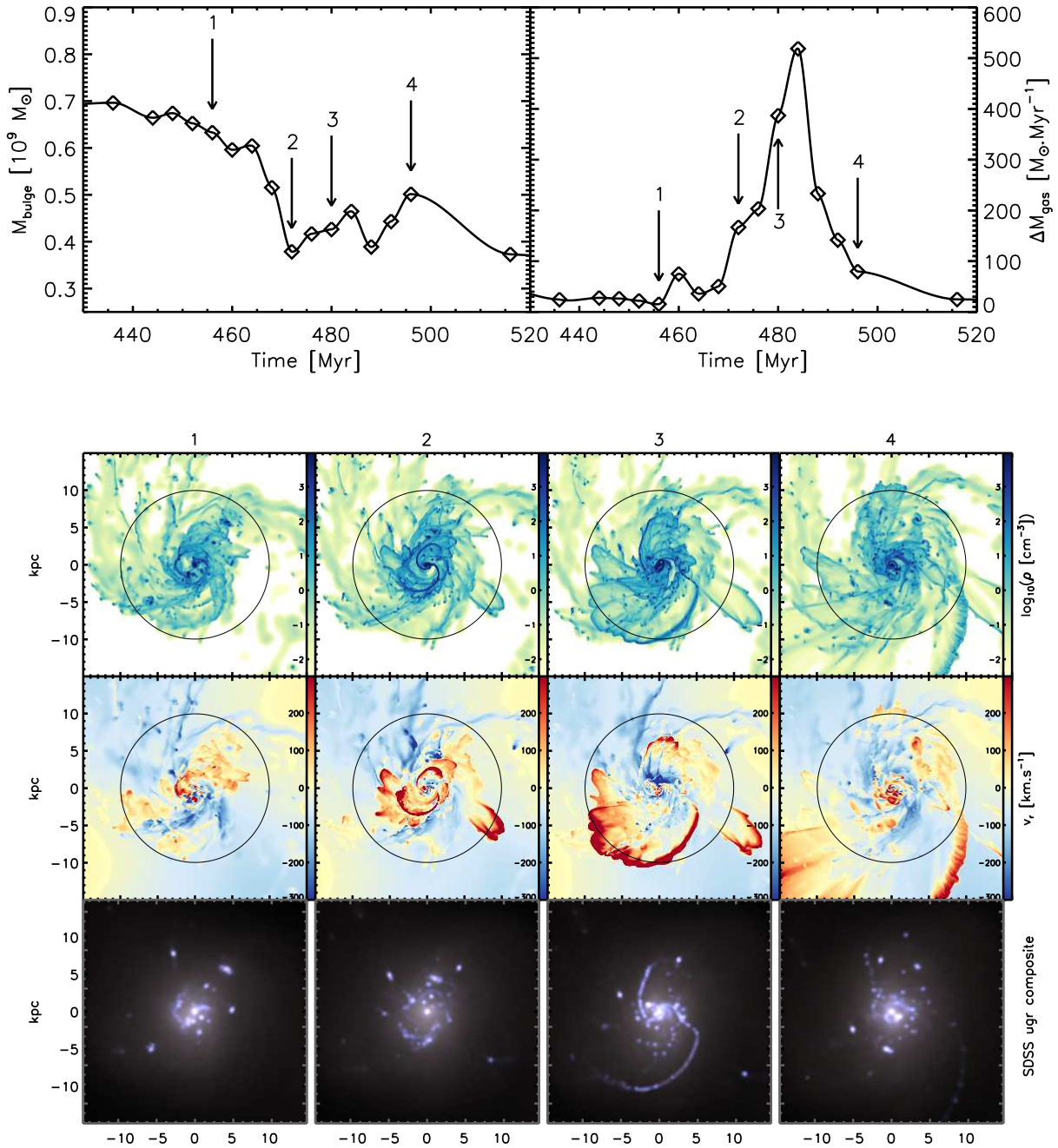


Figure 4.22 – Illustration du processus de coalescence des granules à quatre étapes chronologiques (notées de 1 à 4) de la simulation G1_G1_90_90_180. En haut : Évolution de la masse du bulbe stellaire (à gauche) et du taux d'éjection de masse de gaz (à droite). En bas : Cartes montrant l'évolution de la densité de gaz moyenne le long de la ligne de visée (en haut), de la vitesse radiale cylindrique dans le plan du disque (au milieu), et de l'émission stellaire modélisée avec *STARBURST99* (en bas) à quatre instants distincts repérés sur les courbes d'évolution de la masse du bulbe et du taux d'éjection de gaz. L'émission stellaire est montrée grâce à une image composite des bandes *ugr* à la résolution des observations SDSS pour un décalage spectral $z \sim 0.01$. Le cercle de 10 kpc de rayon sur les cartes de densité et de vitesse radiale du gaz permet de visualiser la sphère dans laquelle est mesuré le taux d'éjection du gaz.

les parties centrales des complexes granulaires les plus massifs ne sont pas en mesure de convertir du gaz en étoiles. En modifiant la loi de formation stellaire, toutes les cellules de gaz à moins de $\Delta T_{poly} = 2 \times 10^5$ K du polytrophe, y compris celles avec des densités $\rho_{gas} > 3500 \text{ cm}^{-3}$, sont autorisées à convertir de la masse gazeuse en particules stellaires. J'ai donc re-simulé trois des simulations de l'échantillon MIRAGE : G1, G2 et G2_G2_90_90_180 (cf. section 3.4). Ces trois nouvelles simulations sont annotées de la lettre "P" (pour prime) dans toutes les figures du chapitre. Les figures 4.23, 4.24, 4.25 présentent les cartes de densité, température, vitesses, et émission stellaire associées à chacune de ces simulations complémentaires. Le choix de ne pas former d'étoiles dans les régions les plus denses dans l'échantillon MIRAGE était motivé par la volonté d'étudier de manière arbitraire le cas des granules survivant à la rétro-action stellaire. En choisissant une limite fixe de température pour la formation stellaire et une énergie injectée par supernova élevée comparée aux valeurs standards, j'ai été en mesure d'obtenir des simulations avec des taux d'éjections de gaz importants (de l'ordre de grandeur du SFR) mais produisant des granules ayant une longue durée de vie (supérieure à un temps orbital).

De manière générale, les simulations complémentaires contiennent beaucoup moins de granules massifs du fait de leur destruction cyclique par les mécanismes de rétro-action stellaire dès le début de la simulation. Le gaz semble être moins turbulent après s'être relaxé depuis les conditions initiales. J'explique la relaxation du gaz par les interactions moins prononcées entre le bulbe stellaire et les granules persistants.

4.4.2 Localisation des complexes granulaires

Afin de comprendre l'origine des mécanismes à même d'éjecter de grandes quantités de gaz par coalescence de granules, j'ai procédé à la caractérisation de ceux-ci dans toutes les simulations de l'échantillon MIRAGE. L'objectif étant d'étudier les propriétés statistiques des granules à chaque pas de temps de 40 Myr et non pas de suivre l'évolution individuelle de ces granules au cours du temps, j'ai utilisé une méthode algorithmique qui permet de détecter les granules sur un instantané. Dans la publication présentée ce chapitre, nous avons au contraire suivi l'évolution des granules dans le temps.

Pour cette méthode algorithmique, j'ai utilisé des cartes de densité du gaz pour une projection de face du disque de chacune des simulations. Les pas de temps des simulations où les deux disques n'ont pas encore fusionné ne sont pas utilisés de manière à éviter des géométries complexes qui rendraient une détection des granules difficile. Chaque carte de densité est composée de 512×512 pixels, le pixel mesurant $120 \times 120 \text{ pc}^2$. Les pixels qui contiennent une valeur de densité moyenne au-dessus de 500 cm^{-3} sont considérés comme faisant partie d'un granule. Afin de déterminer quels sont les pixels faisant partie d'un seul et même granule, j'ai réutilisé l'algorithme d'automate cellulaire 2D présenté dans le chapitre 1 et qui est capable de détecter et classer les groupes de pixels contigus encodant une valeur de densité supérieure à 500 cm^{-3} . Chaque groupe de pixels contigus est considéré comme un granule dans la simulation hydrodynamique 3D. Pour chaque groupe de pixels, je calcule la position du barycentre dans le plan de l'image, donc dans le plan du disque, qui sera considéré comme le barycentre du granule dans le plan du disque. Reste à déterminer la position du granule dans la troisième dimension perpendiculaire au disque. La coordonnée cartésienne verticale z est déterminée en réalisant un histogramme des positions z des cellules AMR de gaz présentes dans un rayon de 100 parsecs autour de la coordonnée (x, y) du granule. Cette méthode a cependant pour défaut de détecter à la fois les granules et les structures en filaments qui ne se sont pas encore fragmentées pour former des granules.

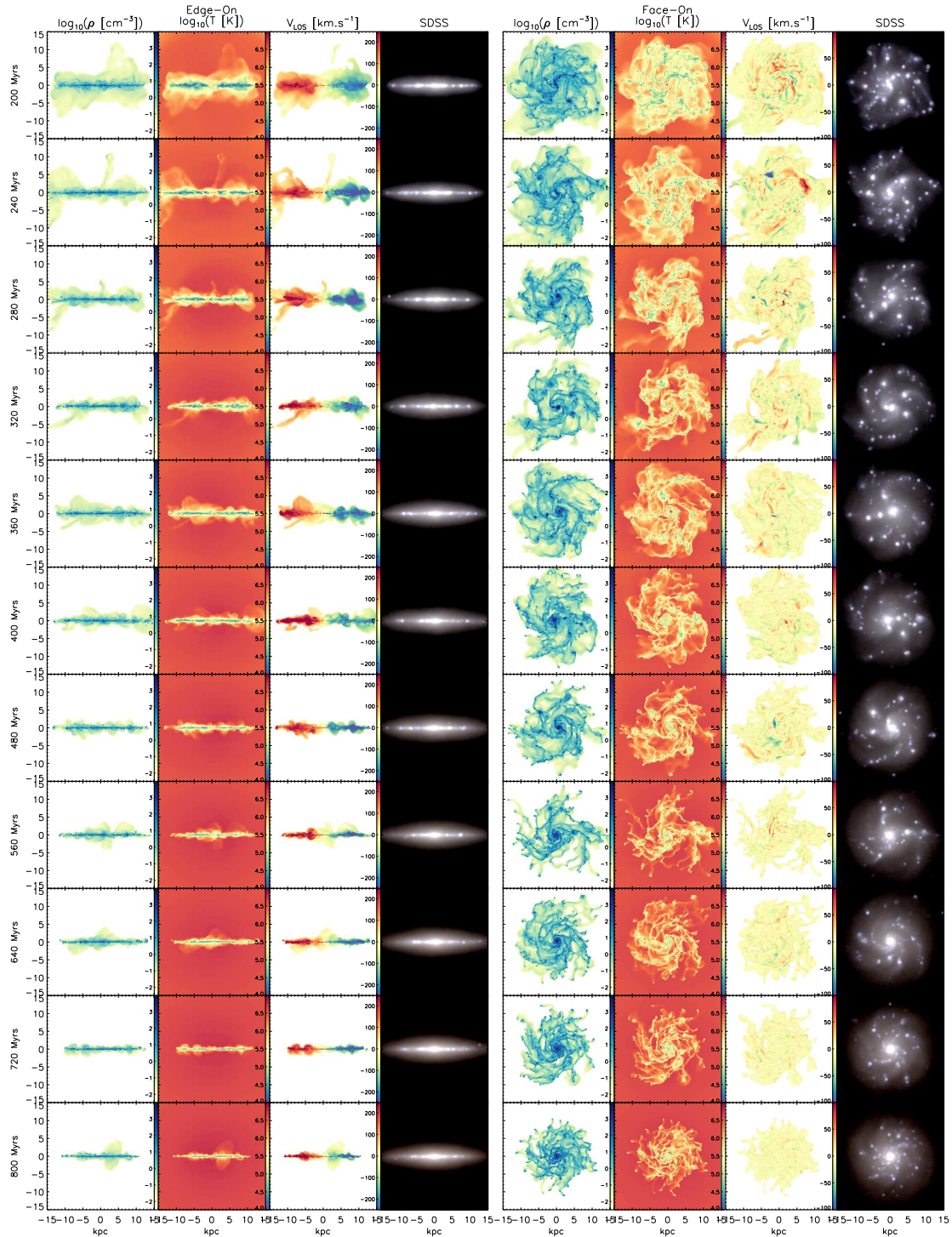


Figure 4.23 – Projection orthographique de la simulation G1P. Deux lignes de visée perpendiculaires sont affichées sur deux colonnes de figures distinctes. La projection de tranche du disque est affichée sur les quatre colonnes de gauche, tandis que la projection de face du disque est affichée sur les quatre colonnes de droite. Les angles de projection sont gardés constants par rapport au référentiel cartésien de la grille AMR de la simulation. Pour chaque projection sont affichées : une carte de densité moyenne pondérée par la masse (première colonne), une carte de température moyenne du gaz pondérée par la masse (deuxième colonne), une carte de vitesses azimutales (i.e. le long de la ligne de visée) et une image composite de pseudo-observations SDSS avec les bandes passantes $u/g/r$ (quatrième colonne). Les échelles de densité, température et vitesses azimutales sont affichées dans la partie intérieure droite de chaque image.

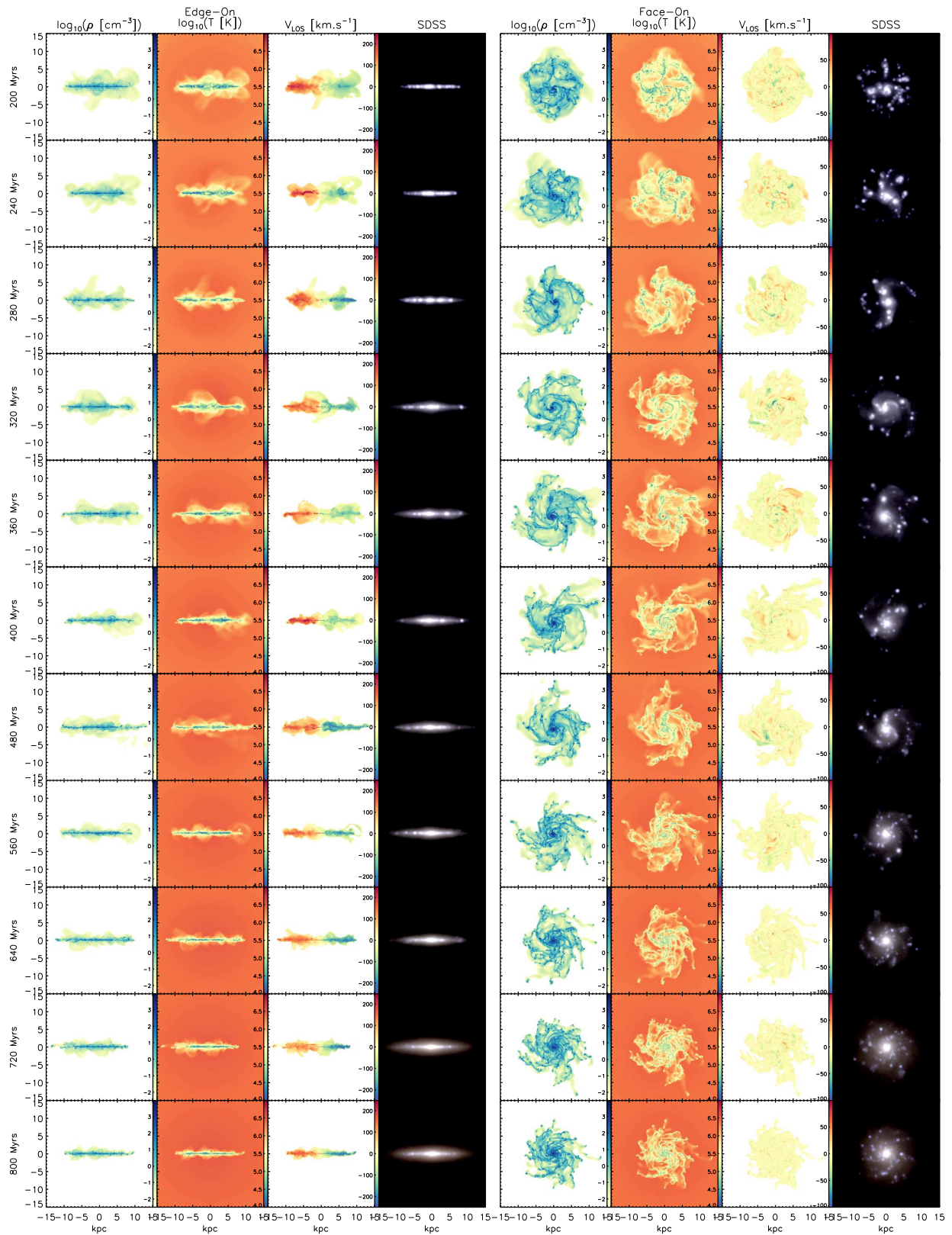


Figure 4.24 – Identique à la figure 4.23, mais pour la simulation G2P.

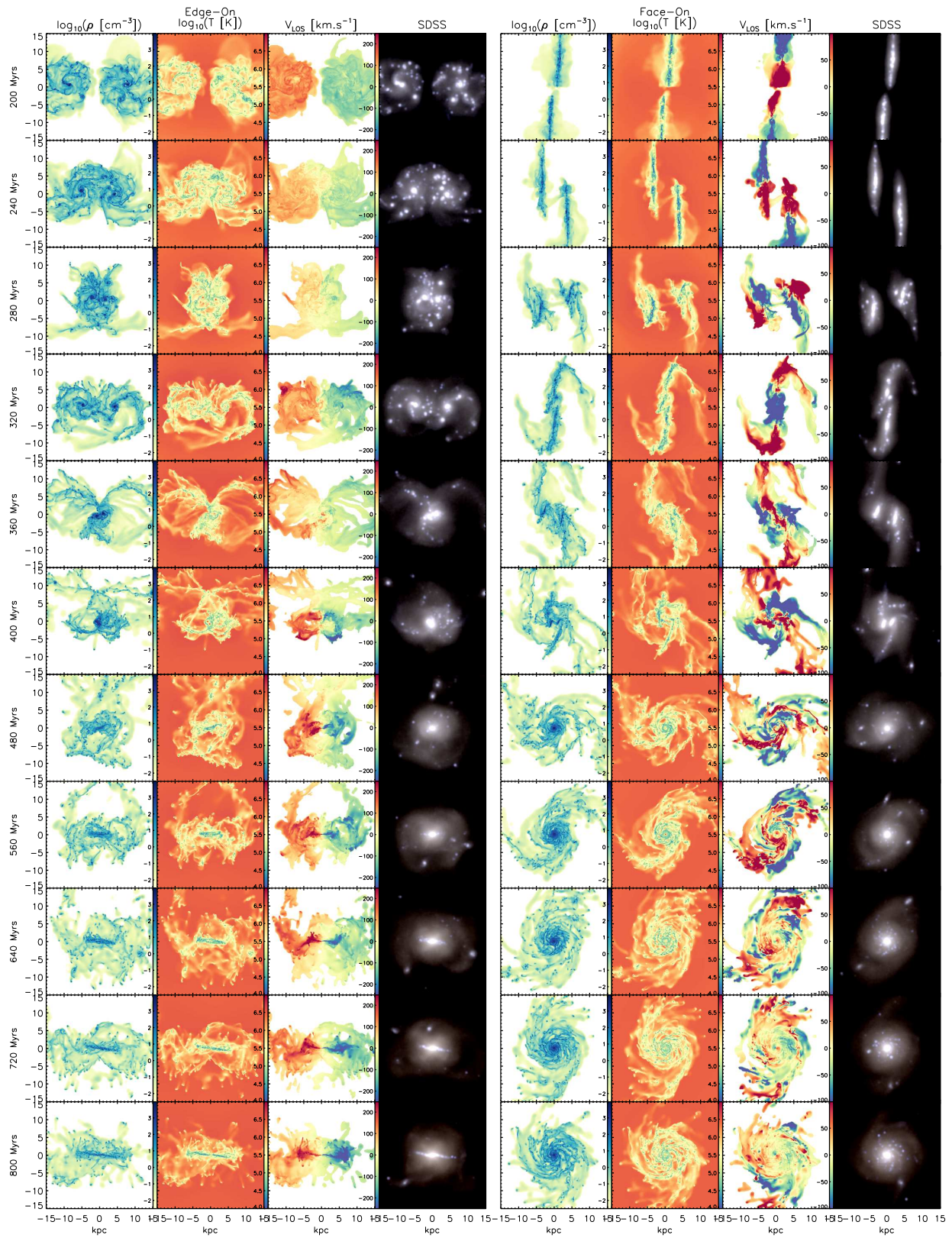


Figure 4.25 – Identique à la figure 4.23, mais pour la simulation G2_G2_90_90_180P.

J'évalue l'aplatissement du groupe de pixels composant le candidat afin de supprimer les filaments qui forment des structures plus allongées que les granules : les groupes de pixels représentant moins de 30% de la surface du carré dans lequel ceux-ci sont inscrits sont considérés comme des filaments. La détection des complexes granulaires est illustrée dans la figure 4.26 pour la simulation G1_G1_90_90_0.

Je supprime ultérieurement ces structures en filament de la liste des granules détectés en évaluant l'aplatissement du groupe de pixels composant le candidat et en supprimant les groupes de pixels les plus aplatis (moins de 30% de pixels dans le carré qui contient tous les pixels détectés). La détection des complexes granulaires est illustrée dans la figure 4.26 pour la simulation G1_G1_90_90_0.

4.4.3 Propriétés statistiques des complexes granulaires

Pour chaque pas de temps de chacune des simulations analysées dans cette étude, nous disposons des informations de position, de taille, de masse et de vitesse de chacun des granules. La propriété physique la plus importante qui semble modifier le mécanisme de coalescence des granules est la fraction de gaz de ces derniers. Le pourcentage de masse baryonique des granules en fonction de leur fraction de gaz f_g est tracée dans la figure 4.27.

Pour chaque simulation (MIRAGE ou P), les granules sont classés en fonction de leur fraction de gaz (de 0 à 1, avec un pas de 0.02) quel que soit le pas de temps de la simulation, l'objectif n'étant pas ici de suivre l'évolution des granules en fonction du temps mais la répartition de la masse baryonique en fonction du contenu en gaz de ces granules. Cette figure montre clairement que la majorité de la masse se trouve à des concentrations très élevées en gaz ($f_g > 0.8$) pour les simulations MIRAGE. Ce constat diffère pour les simulations P, dont la majorité de la masse des clumps se trouve à des fractions de gaz plus faibles ($f_g < 0.4$). Cette différence s'explique par le taux de formation stellaire beaucoup plus élevé dans le noyau des granules des simulations P qui transforme cycliquement le gaz en étoiles. La structure stellaire semble être en mesure de persister, de telle manière que le gaz s'effondre un peu plus tard dans le centre du granule stellaire rémanent.

Afin de comprendre comment les granules accumulent et maintiennent une fraction de gaz aussi élevée, je montre la carte de métallicité pour la simulation G1P dans la figure 4.28, avec des paramètres de projection identiques à ceux de la figure 4.26. La position des granules est indiquée par les symboles en forme de losange, ce qui permet de visualiser l'environnement direct des granules. De nombreux granules, notamment dans les régions externes du disque, semblent alimentés par du gaz à faible métallicité et qui provient donc du milieu circum-galactique. Il est frappant de constater la faible corrélation entre la métallicité et la densité des granules, ces dernières se trouvant indistinctement dans des régions de faible ou de forte métallicité, en particulier dans les régions externes des galaxies et hors du plan équatorial de la galaxie ou l'on observe des régions de faible densité très métalliques qui ne contiennent pas de granules. Afin d'être certain que ce gaz soit chaud et avec une faible métallicité soit bien en train de s'effondrer sur les granules, j'ai aussi réalisé des cartes des composantes cylindriques radiales et verticales de la vitesse du gaz dans le disque (figure 4.29) et des cartes de température du gaz dans le disque (figure 4.30). La vue de face du disque de la figure 4.29 permet de visualiser le gaz qui s'effondre radialement vers le centre (vitesses négatives), ainsi que le gaz en train de s'éloigner du centre de masse (vitesses positives), on y observe des mouvements antagonistes d'éloignement du centre (en rouge) et de chute vers le centre (en bleu) tous deux dans le plan du disque. On observe

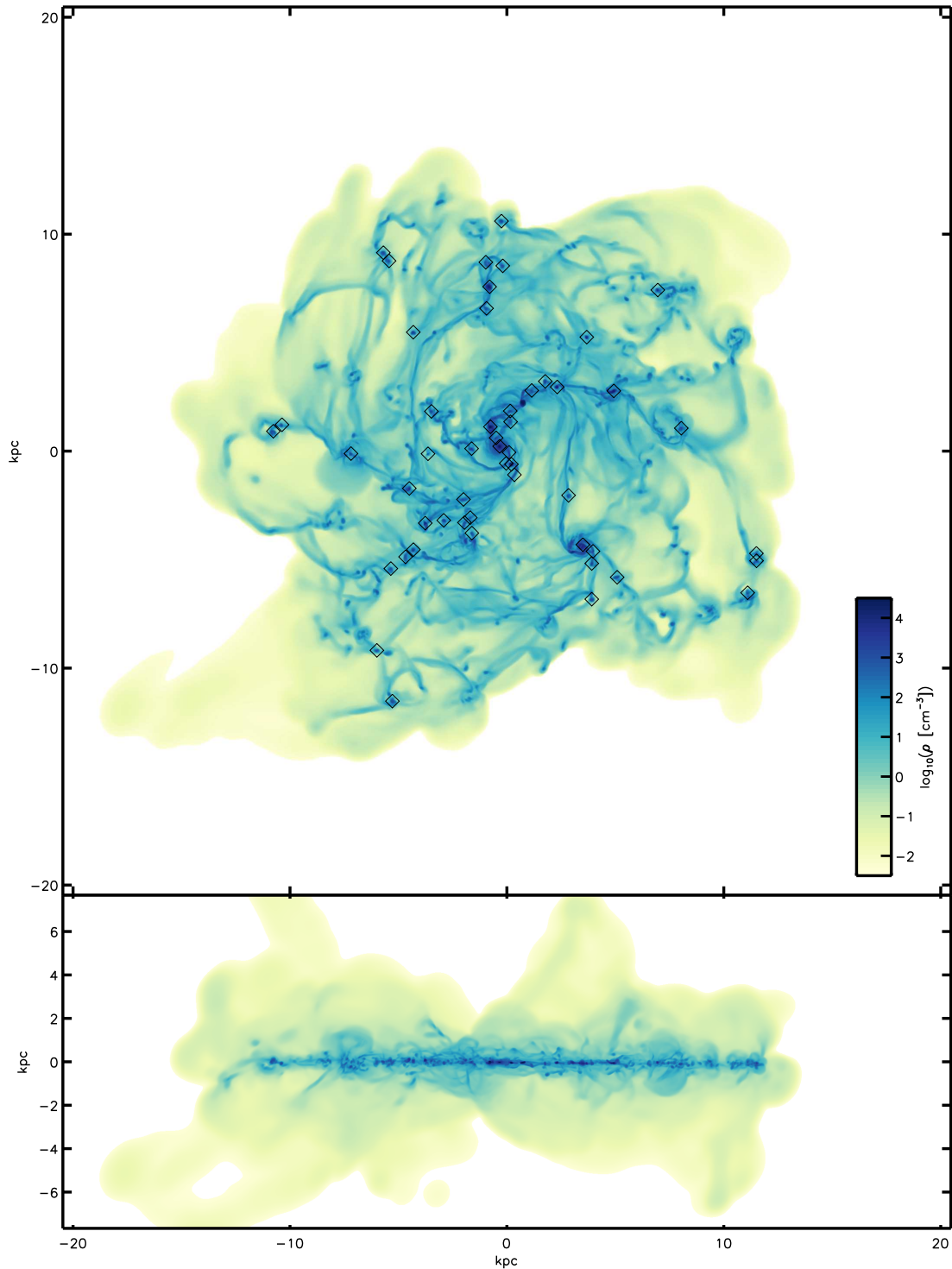


Figure 4.26 – Vue de face (figure du haut) et vue de profil (figure du bas) de la carte de densité moyenne pondérée par la masse des cellules AMR le long de la ligne de visée pour la simulation G1P à 200 Myr après le lancement de la simulation. Chacun des symboles "◇" note la position des granules de densité supérieure à 500 cm^{-3} identifiés par l'algorithme de détection.

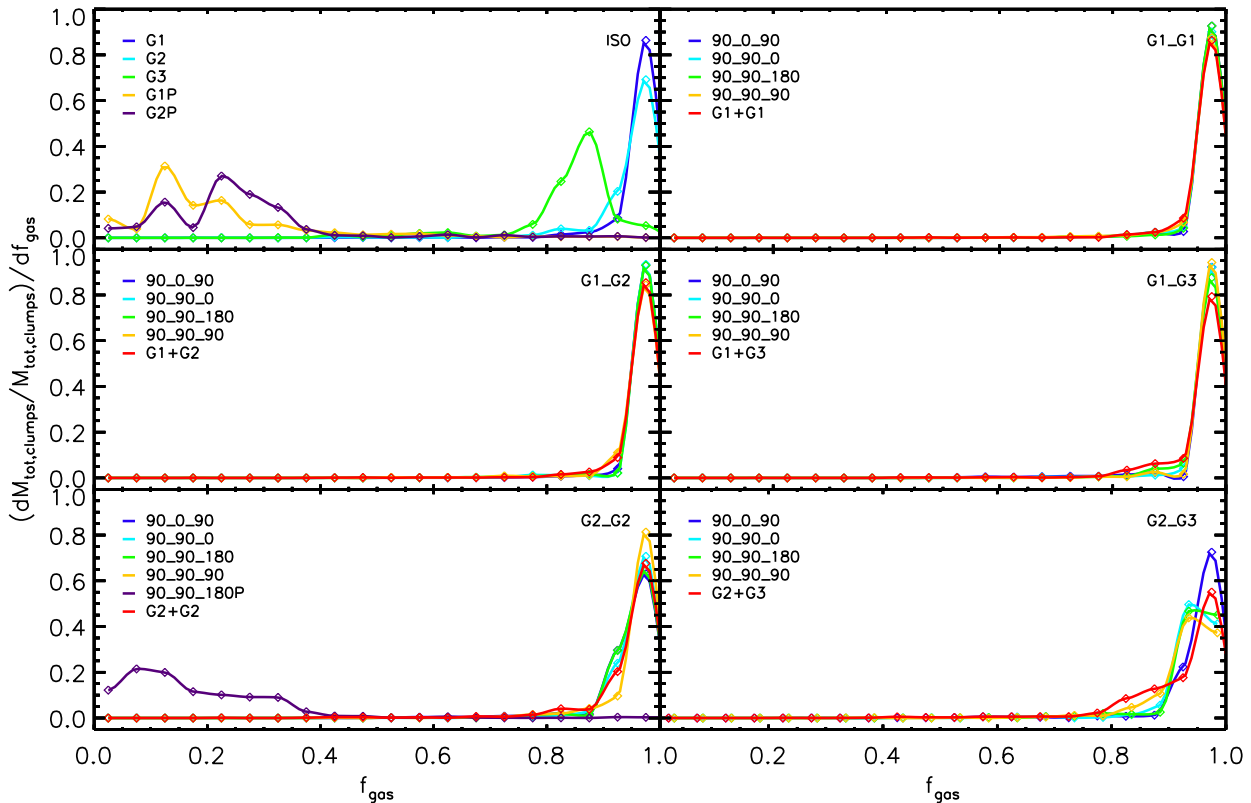


Figure 4.27 – Pourcentage de masse baryonique contenue dans les complexes granulaires en fonction de la fraction de gaz des granules pour l’ensemble des simulations MIRAGE et pour les simulations complémentaires P. Pour chaque simulation, les granules sont triés en fonction de leur fraction de gaz, indépendamment du pas de temps. La masse baryonique totale des granules, ayant une fraction de gaz donnée, est calculée en effectuant la somme de leur masse baryonique, quel que soit le pas de temps d’où le granule est extrait. Cette masse par unité de fraction de gaz est transformée en pourcentage en normalisant par la masse baryonique totale de l’ensemble des granules, toute fraction de gaz confondue. Chaque panneau explore les orientations de disques pour des masses de galaxies initiales fixes. Sur chaque panneau qui concerne une configuration de fusion, la courbe rouge montre la somme des distributions des deux disques qui ont évolué de manière isolée. L’essentiel de la masse des granules est comprise dans les granules très gazeuses pour les simulations MIRAGE alors qu’elle est contenue dans les granules peu gazeuses pour les simulations P.

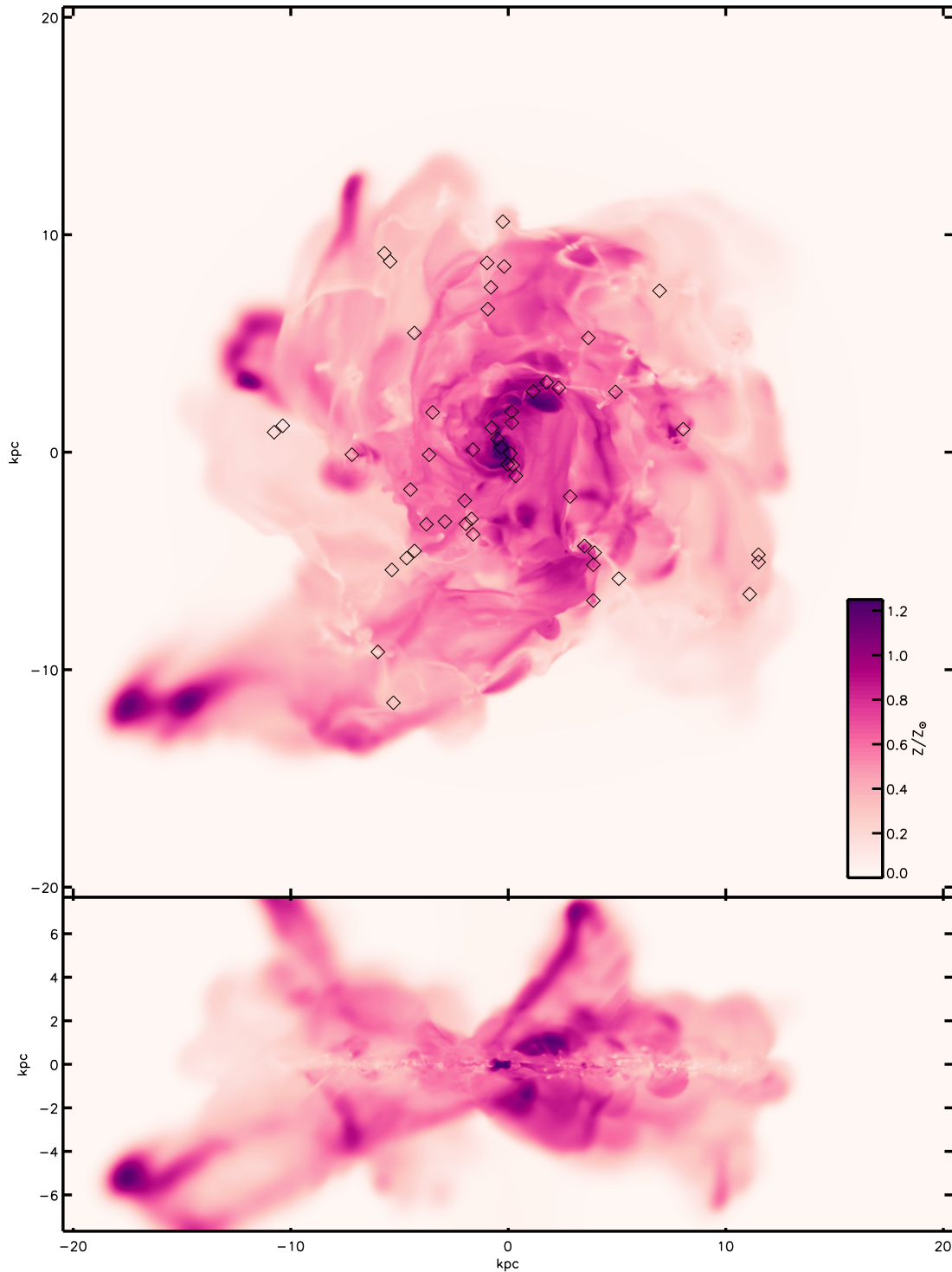


Figure 4.28 – Vue de face (figure du haut) et vue de profil (figure du bas) de la carte de métallicité moyenne (en unité de métallicité solaire) pondérée par la masse des cellules AMR le long de la ligne de visée pour la simulation G1P à 200 Myr après le lancement de la simulation. Chacun des symboles "◇" note la position des granules de densité supérieure à 500 cm^{-3} identifiés par l'algorithme de détection.

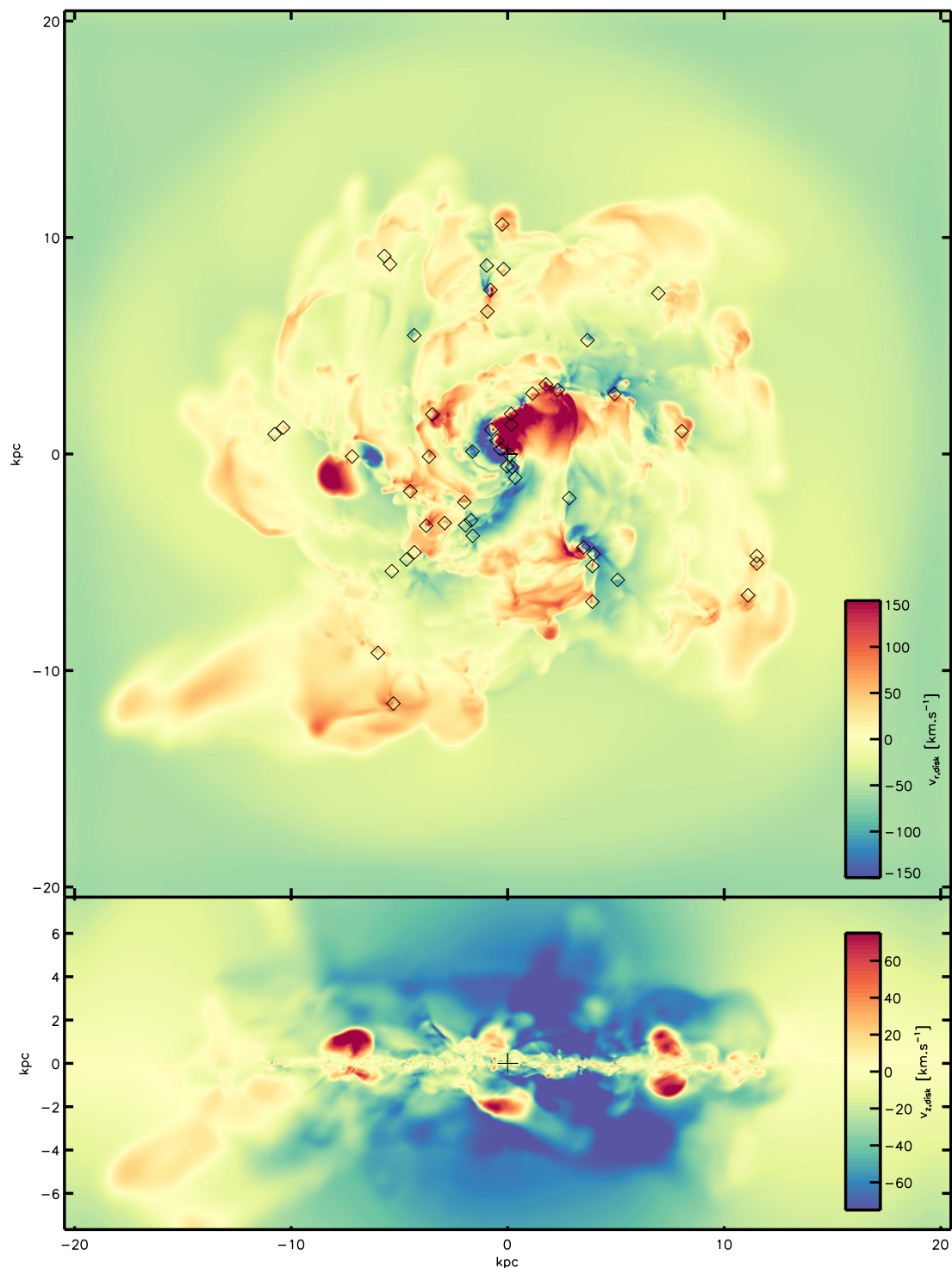


Figure 4.29 – Vue de face (figure du haut) et vue de profil (figure du bas) de la valeur moyenne de la composante de vitesse cylindrique radiale (en haut) et de la composante de vitesse perpendiculaire d'éloignement au plan du disque $v_z z/|z|$ (en bas) pour la simulation G1P à 200 Myr après le lancement de la simulation. Chacun des symboles "◇" note la position des granules de densité supérieure à 500 cm^{-3} identifiés par l'algorithme de détection. Le centre du système de coordonnées déterminé par la procédure de recentrage sur le buble est indiqué par une croix.

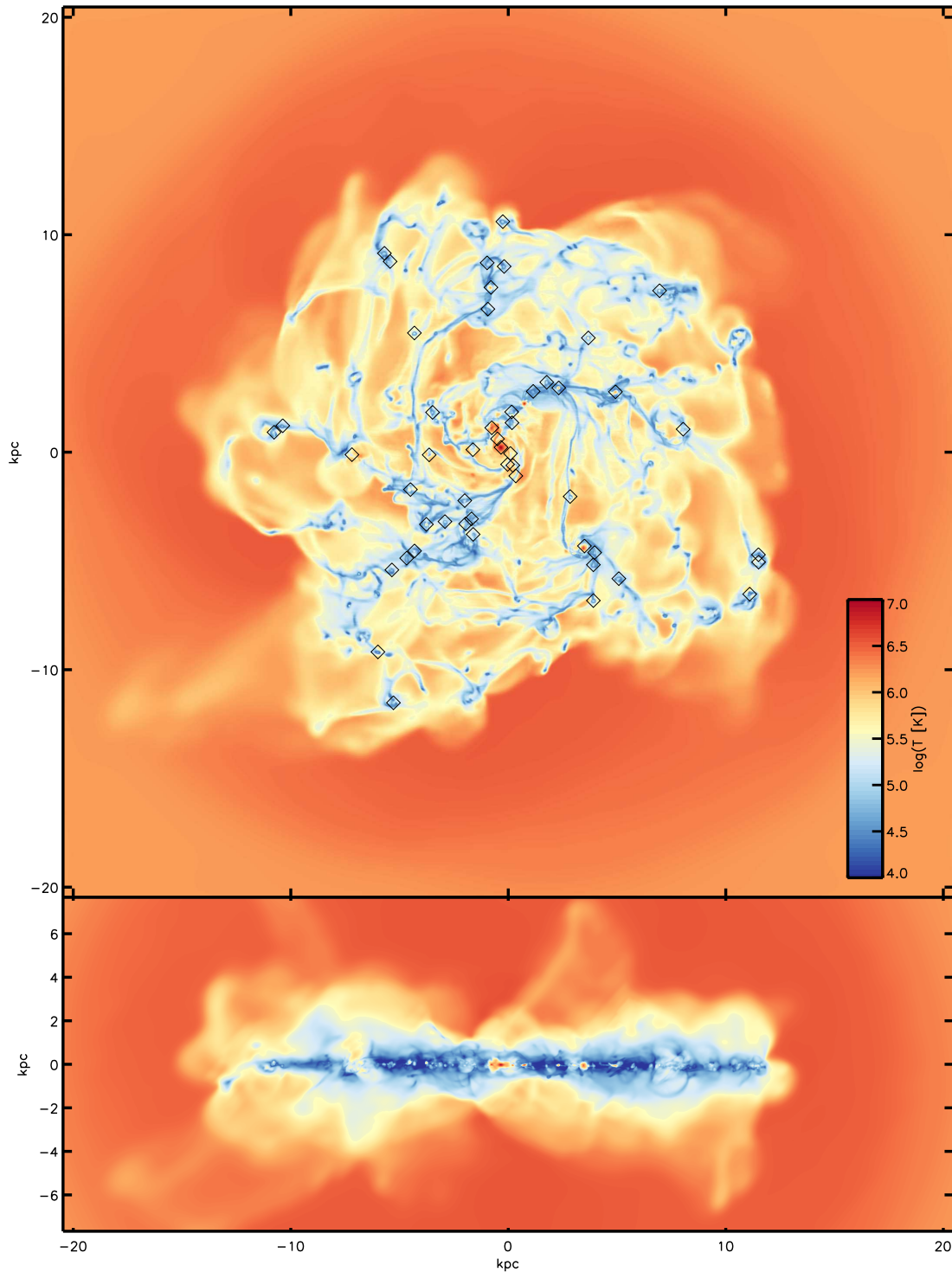


Figure 4.30 – Vue de face (figure du haut) et vue de profil (figure du bas) de la carte de température moyenne (en K) pondérée par la masse des cellules AMR le long de la ligne de visée pour la simulation G1P à 200 Myr après le lancement de la simulation. Chacun des symboles "◇" note la position des granules de densité supérieure à 500 cm^{-3} identifiés par l'algorithme de détection.

dans d'autres régions du disque des comportements cinématiques similaires, toujours associés à la présence de granules. La vue de profil du disque montre la valeur pondérée par la masse la composante de vitesses perpendiculaire au disque $v_z z/|z|$ (i.e. vitesse d'éloignement du disque) : cette carte permet de visualiser le gaz du milieu circum-galactique en train de tomber sur le disque (vitesses négatives en bleu), ainsi que les éjections de gaz en partance du plan du disque provoquées par la rétro-action stellaire (vitesses positives en rouge). On constate sur les cartes de température du disque de gaz vu par la tranche, en provenance du milieu circum-galactique, qui s'effondre sur le disque est chaud ($T > 5 \times 10^6 \text{K}$) et que les courants filamentaires de gaz qui alimentent les granules sont au contraire plutôt froids et peu métalliques. Une inspection de ces trois cartes permet de déduire que (i) le disque est alimenté par l'accrétion de gaz chaud halo circum-galactique de manière isotrope (verticalement et horizontalement) (ii) des courants de gaz pauvres en métaux (donc provenant du milieu intergalactique) existent dans le plan du disque et semblent alimenter les granules en gaz primordial au-delà d'un rayon d'environ 75% de la longueur d'échelle du disque stellaire dans le cas présenté (avec une longueur d'échelle pour le disque stellaire $r_{stars} = 3.9 \text{ kpc}$ dans le cas de G1P).

4.4.4 Éjection de gaz

Comme nous l'avons déjà constaté dans l'article présenté dans ce chapitre, la migration des granules vers le centre de masse de la galaxie par friction dynamique convoie une quantité importante de gaz. Dans le cas de figure où aucun bulbe ne s'est encore formé au centre de masse de la galaxie, les granules sont en orbite quasi-circulaire et peuvent être capturés par leurs voisins les plus massifs donnant rapidement naissance à des granules plus massifs vers le centre de masse. Tous les granules qui sont sur des orbites plus lointaines vont peu à peu migrer vers ce proto-bulbe. L'approche finale d'un granule vers le proto-bulbe se traduit par une forte accélération des granules, dont la vitesse circulaire atteint le pic de la courbe de rotation. La phase gazeuse des granules possède un support dynamique en partie assuré par de la rotation et chacun de ces granules transporte un moment angulaire associé à cette rotation. Du fait que ces granules naissent par cisaillement et effondrement gravitationnel dans le plan du disque, leur moment angulaire est toujours quasi-colinéaire au moment angulaire total du disque (Ceverino et al., 2012). La coalescence des granules est donc un phénomène tout à fait comparable à la fusion de deux galaxies spirales coplanaires à leur plan orbital. Cette configuration assure une forte résonance gravitationnelle : la force de marée subie par les granules tend à les déformer et une fraction conséquente du gaz se retrouve éjectée dans des queues de marée internes au disque (voir les simulations progrades coplanaires de Chilingarian et al. 2010). Ce phénomène est d'autant plus efficace que les granules sont massifs. L'arrivée successive de granules sur le proto-bulbe assure une éjection de grandes quantités de gaz tout au long des 800 Myr simulés dans les simulations MIRAGE. La figure 4.31 présente le taux d'éjection de gaz, mesuré dans une sphère de rayon 20 kpc. Le gaz dont la métallicité est plus grande que $Z > 10^{-3} Z_{\odot}$ (et qui a donc été enrichi par les métaux relâchés par des supernova) et qui est capable de sortir de ce volume sphérique en moins de 5 Myr est comptabilisé pour le calcul du taux d'éjection. Une grande partie du gaz éjecté atteint des vitesses supra-Viriel, i.e. une vitesse supérieure à 300 km.s^{-1} et sera amenée à échapper au potentiel gravitationnel de la galaxie. Les éjectas les moins rapides sont soumis à un ralentissement, notamment par viscosité hydrodynamique du fait de la densité non-négligeable du milieu circum-galactique ($\sim 10^{-3} \text{ cm}^{-3}$). Au total, les simulations G1_G1 sont capables d'éjecter environ 20% de leur masse initiale de gaz. Ce

mécanisme est d'autant plus prononcé que la galaxie est massive. En effet, les galaxies les moins massives abritent des granules peu massifs qui sont au contraire susceptibles d'être plus endommagés par la rétro-action stellaire (voir la distribution des fractions de gaz de la simulation G3 sur la figure 4.27).

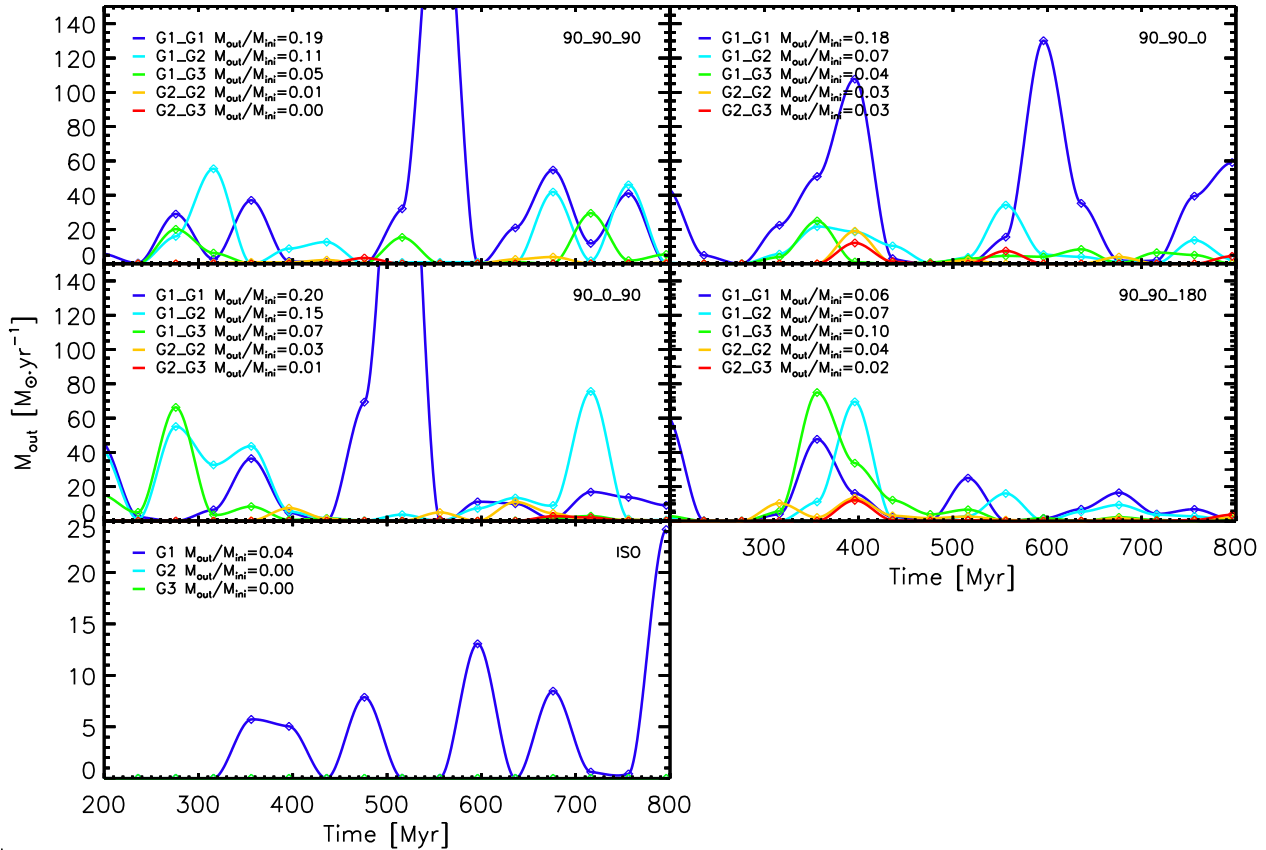


Figure 4.31 – Évolution du taux d'éjection de gaz au cours du temps pour chacune des simulations de l'échantillon MIRAGE. Pour chaque pas de temps, on évalue la masse de gaz capable de sortir d'une sphère de rayon 20 kpc centrée sur le centre cinématique du système en moins de 5 Myr. Seul le gaz plus métallique que $Z > 10^{-3} Z_{\odot}$ est considéré pour évaluer le taux d'éjection de masse de gaz. Sur chaque panneau est indiquée la quantité totale de masse M_{out} ayant quitté la galaxie en unité de masse initiale de gaz M_{ini} .

Bien qu'il y ait une quantité plus importante de supernova dans les simulations P que dans les simulations MIRAGE, je n'observe aucune éjection capable de s'éloigner de plus de 20 kpc du disque. En effet, l'accrétion de gaz du milieu circum-galactique semble être très efficace pour ralentir le gaz expulsé par les explosions de supernova. En effet, ces éjections sont bien moins collimatées que celles provoquées par la fusion de granules : elles s'étalent sur une surface plus grande, ce qui permet au gaz en chute sur le disque de ralentir par friction hydrodynamique les baryons lancés par les supernova. La destruction cyclique des granules dans les trois simulations P, même les plus massifs ($\sim 10^9 M_{\odot}$) diminue la fraction de gaz moyenne de ces granules et par conséquent rend leur fusion inefficace pour éjecter du gaz au-delà des régions les plus centrales.

4.4.5 Régulation du bulbe

La coalescence de granules résultant d'une instabilité gravitationnelle affecte la distribution spatiale de gaz mais affecte aussi celle des étoiles, plus particulièrement celle du bulbe. L'interaction de granules très riches en gaz est capable d'arracher une partie des étoiles déjà présentes dans le puits de potentiel central de manière si efficace que le bulbe perd constamment de la masse dans la majorité des simulations MIRAGE (voir figures 4.32 et 4.33). Les bulbes stellaires dans le jeu de simulations complémentaires P voient quant à eux leur masse augmenter pour atteindre un niveau supérieur à $10^{10} M_{\odot}$. Les fusions de granules dans l'échantillon MIRAGE ne permettent pas la création de bulbes plus massifs que $2.5 \times 10^9 M_{\odot}$.

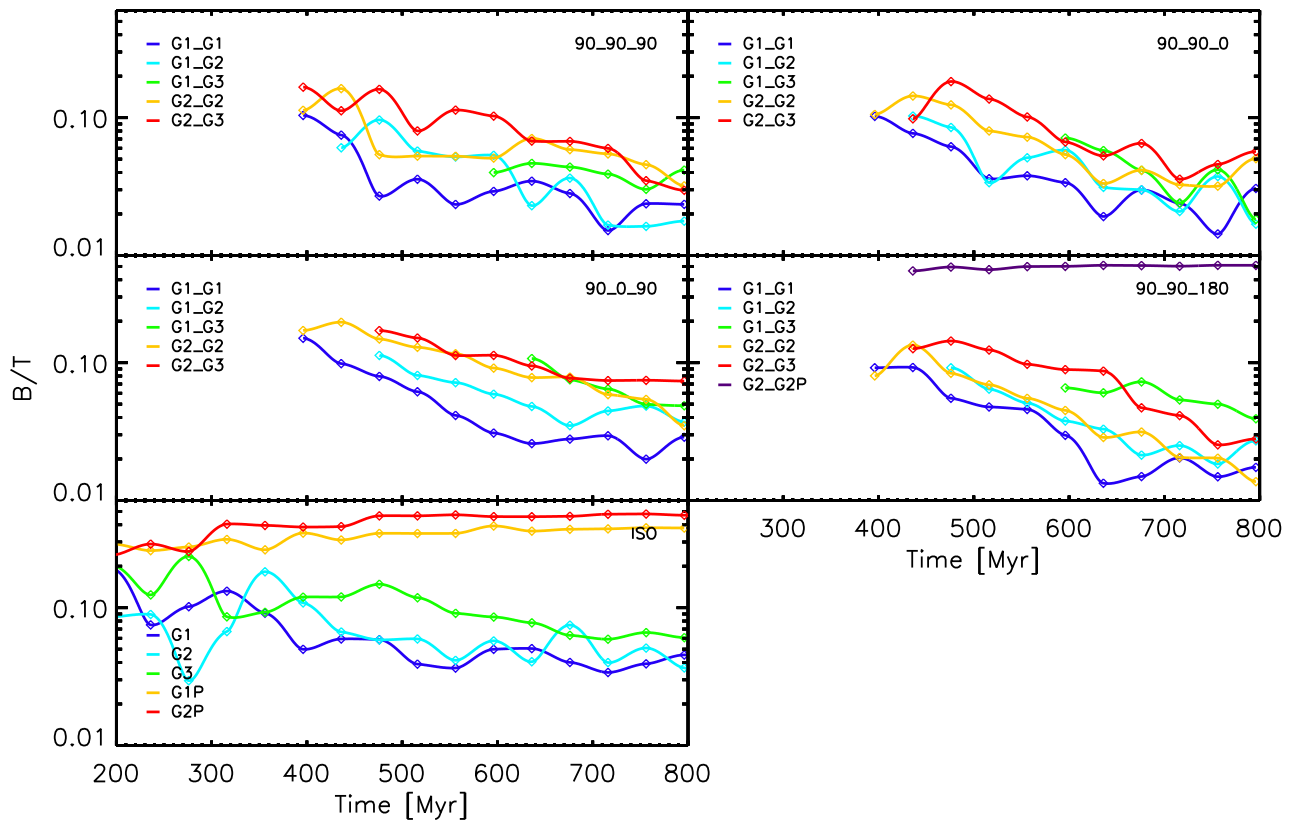


Figure 4.32 – Évolution temporelle du rapport de masse du bulbe sur la masse stellaire totale. Les masses sont définies par l'ajustement d'un profil exponentiel pour le disque que l'on soustrait au profil mesuré pour obtenir la masse du bulbe stellaire. Les mesures sont réalisées chaque 40 Myr à partir de la coalescence des deux galaxies (~ 400 Myr pour les fusions les plus rapides). Afin d'accroître la clarté de la figure, les mesures d'une même simulation sont reliées entre elles par une interpolation cubique. Chaque panneau explore une orientation initiale de disques (G_i-G_j) indiquée par $\theta_1-\theta_2-\kappa$. Le rapport B/T est de 8% dans les conditions initiales de toutes les simulations.

Une inspection visuelle comparative entre simulations MIRAGE et simulations P sur des images composites trois bandes (filtres u, g et r) construites à partir de pseudo-observations SDSS (figure 4.34) permet d'observer un bulbe stellaire beaucoup plus proéminent dans le cas où la formation stellaire est non-bridée en température. Les disques des simulations G1 et G2 sont sensiblement plus épais que ceux des simulations G1P et G2P du fait d'une

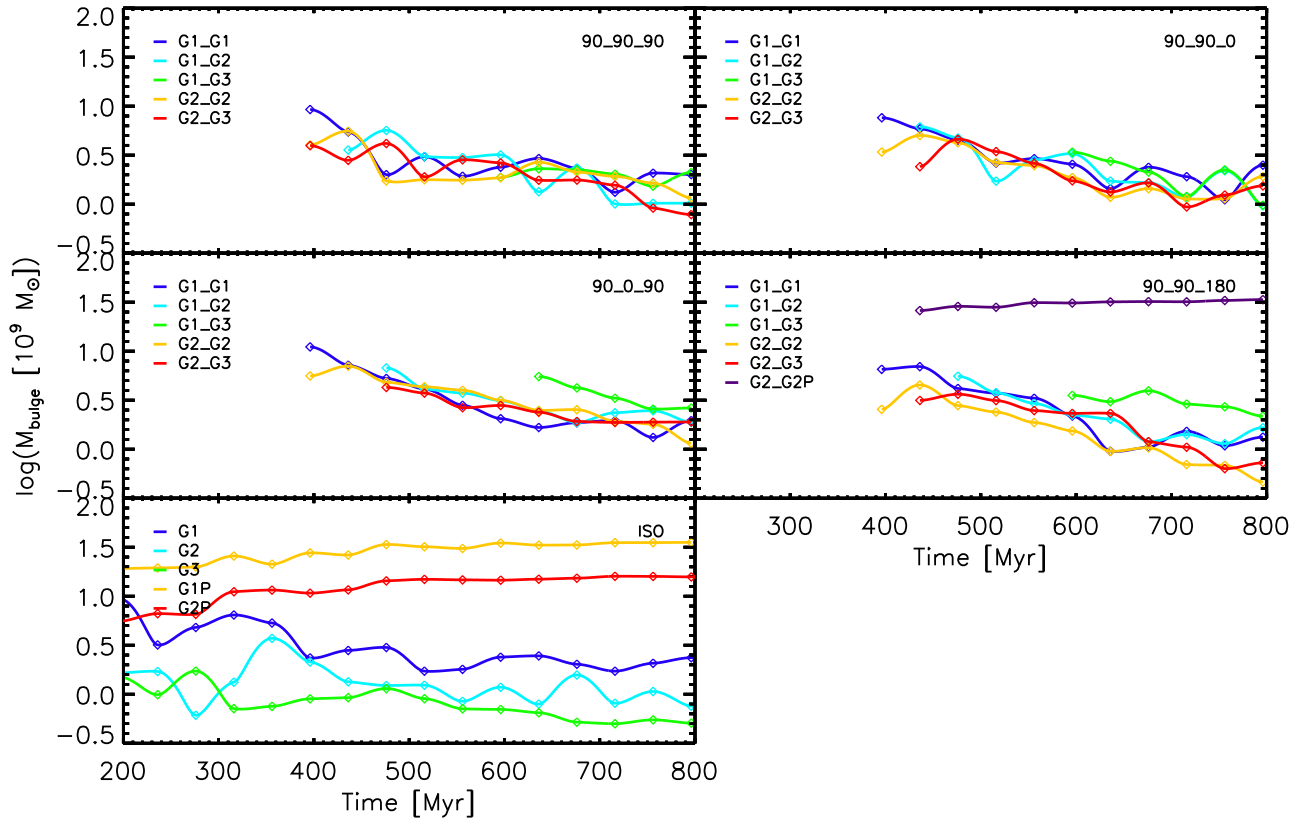


Figure 4.33 – Évolution temporelle de la masse stellaire du bulbe. Les mesures sont réalisées chaque 40 Myr à partir de la coalescence des galaxies. Afin d'accroître la clarté de la figure, les mesures d'une simulation sont reliées entre elles par une interpolation cubique. Chaque panneau explore une orientation initiale de disques (G_i-G_j) indiquée par $\theta_1-\theta_2-\kappa$.

redistribution des étoiles sur des orbites plus excentriques qui vont venir peupler un ovoïde stellaire. La simulation de fusion G2_G2_90_90_180P présente des reliques de queues de marées, tandis que la simulation G2_G2_90_90_180 en est exempte. Le flux stellaire apparaît sensiblement plus concentré dans la simulation G2_G2_90_90_180P que dans son analogue de l'échantillon MIRAGE. En effet, les fusions de galaxies sont des processus hautement dissipatifs pour le gaz et permettent de concentrer la matière baryonique dans un volume très restreint (Bournaud et al., 2011). Les fusions successives de granules riches en gaz dans la simulation G2_G2_90_90_180 permettent d'éviter cet effet en redistribuant rapidement les baryons dans le disque (pour le gaz) et dans un ovoïde (pour les étoiles).

Afin de cerner la nature du processus physique à l'origine de l'expulsion des étoiles du bulbe, j'ai identifié dans chacune des simulations les particules stellaires gravitationnellement liées au bulbe. Pour ce faire, je sélectionne les particules à l'intérieur d'un rayon $r_{\text{limit}}=0.3r_{\text{stars}}$ soit 30% de la longueur d'échelle du disque. Afin de s'assurer que ces particules sont gravitationnellement liées au potentiel du central du bulbe, je sélectionne celles dont la vitesse est inférieure à 50% de la vitesse d'échappement locale $v_{\text{esc}} = \sqrt{2GM_{\text{cumul}}(r)/r}$ avec $M_{\text{cumul}}(r)$ la masse cumulée au rayon r . Pour chacune de ces particules, je dispose d'un identifiant unique permettant de retrouver leurs positions à n'importe quel pas de temps de la simulation, ce qui me permet de suivre l'évolution du profil de densité radial ρ_{id} de ces

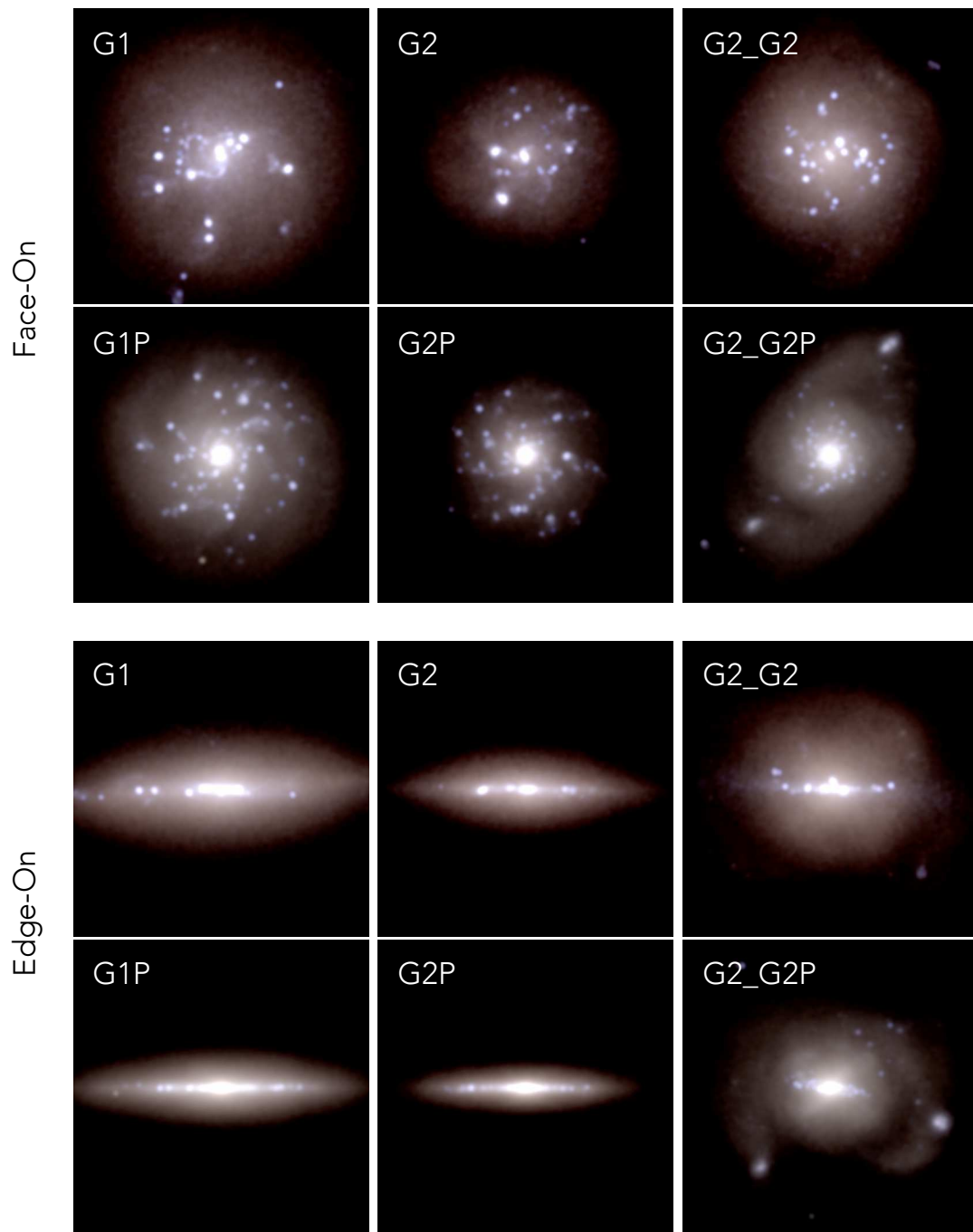


Figure 4.34 – Comparaison d’images composites réalisées à partir de pseudo-observations SDSS dans les bandes passantes $u/g/r$ pour une projection à $z=0.01$ et une résolution spatiale de 1.4 secondes d’arc. Toutes les pseudo-observations sont construites à $t=800$ Myr après les conditions initiales. Le panneau du haut présente des projections où les disques sont vus de face et le panneau du bas présente des projections où les disques sont vus par la tranche. Les références inscrites en haut à gauche des images correspondent aux différentes simulations (cf. section 3.4 et 4.4.1).

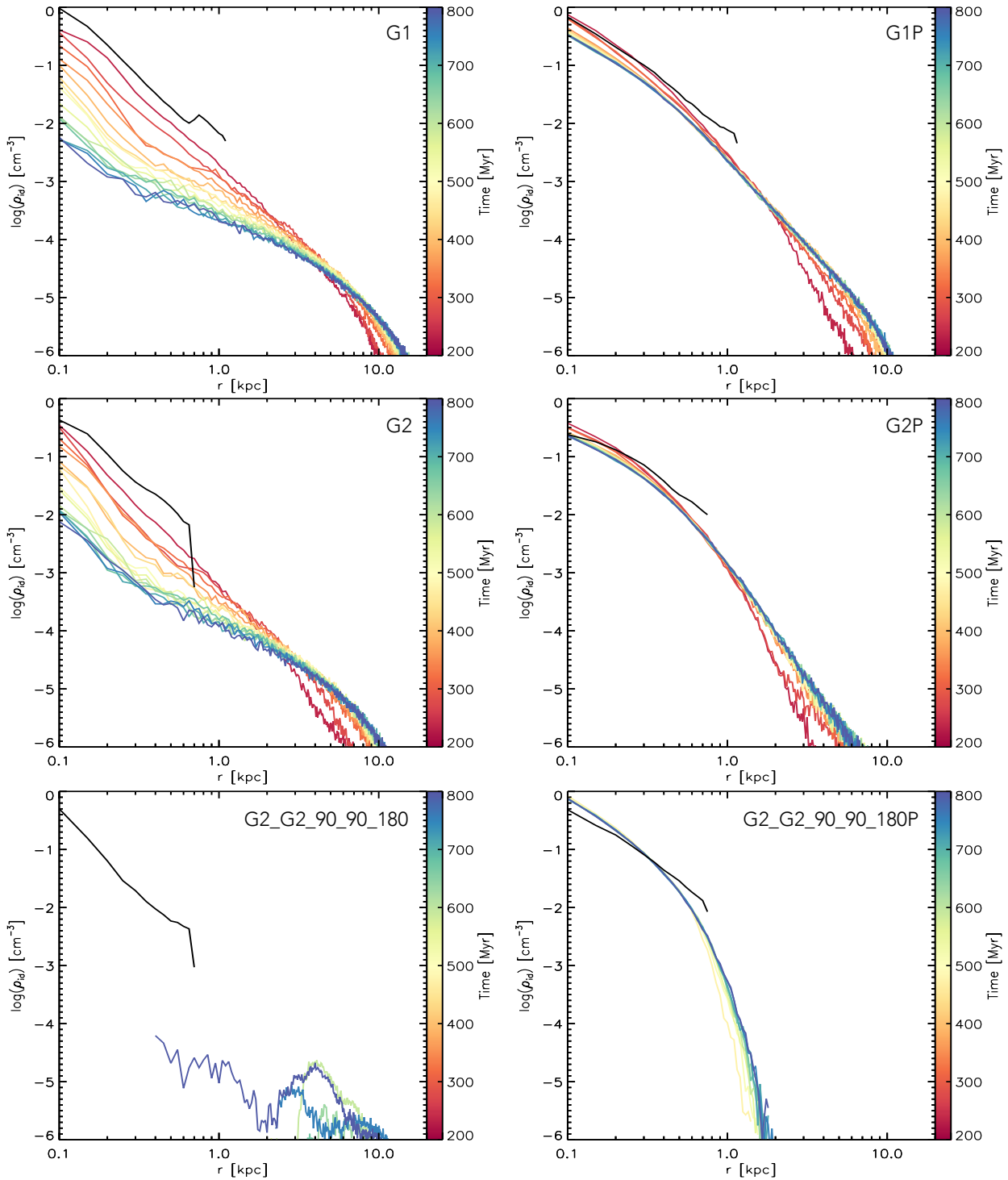


Figure 4.35 – Migration des particules stellaires du bulbe induite par la coalescence de complexes granulaires vers le halo. Un profil initial de densité radiale pour les particules stellaires faisant partie du bulbe est sélectionné à temps t_i , représentée par la courbe noire sur les 6 figures. Pour les deux lignes du haut, le temps initial t_i est choisi à 200 Myr alors que le temps initial dans le cas de la fusion entre deux galaxies (ligne du bas) est fixé au moment de la coalescence ($t_i=380$ Myr dans le cas MIRAGE, colonne de gauche ; $t_i=440$ Myr dans le cas P, colonne de gauche). L'évolution temporelle de la distribution radiale de ces particules est ensuite suivie au cours de chaque simulation grâce à l'identifiant unique associé à chaque particule. Le pas de temps est représenté par la couleur de la courbe de densité. Les identifiants des simulations sont donnés en haut à droite de chaque figure.

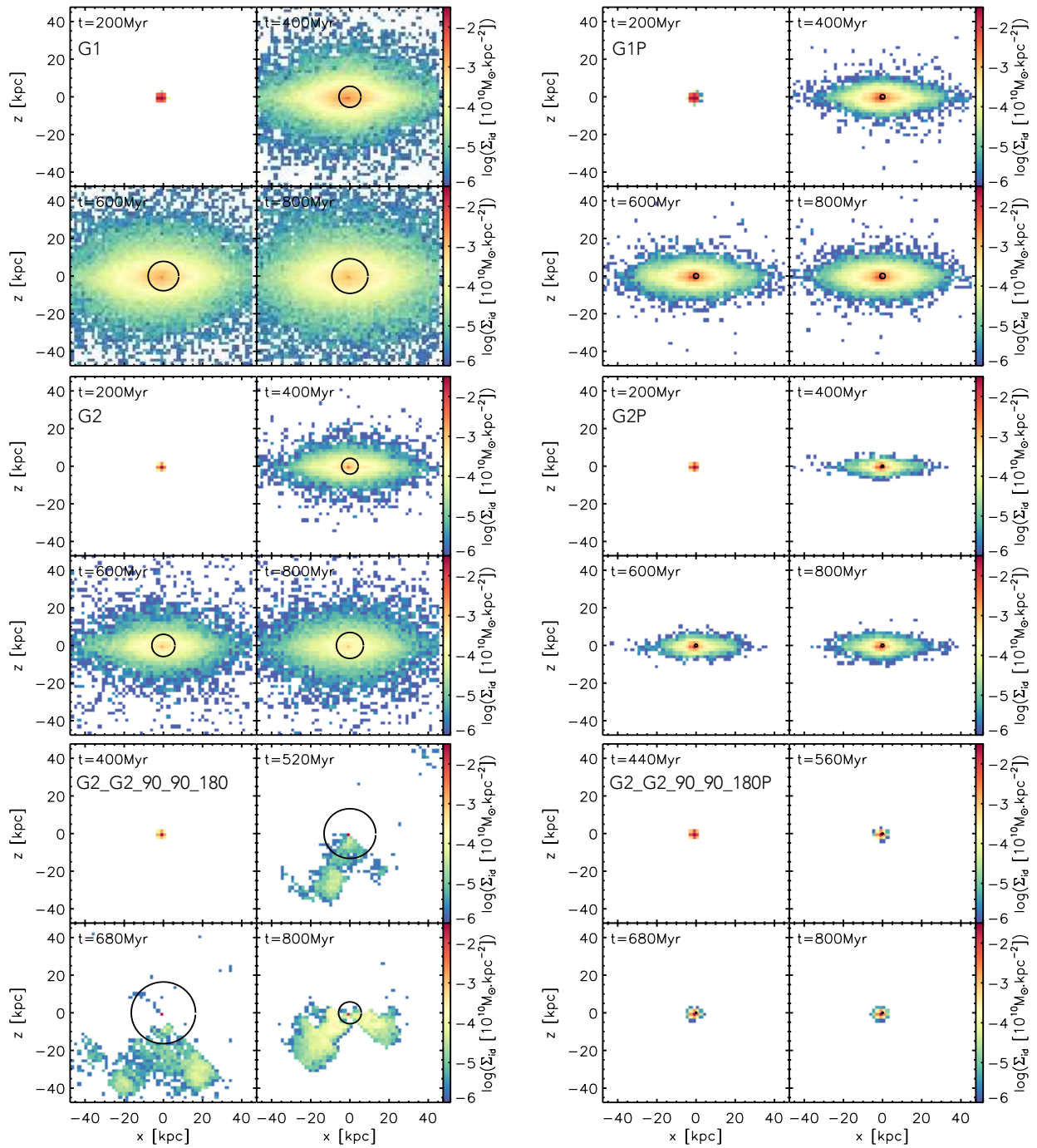


Figure 4.36 – Évolution de la densité surfacique des particules gravitationnellement liées au bulbe stellaire dans un plan orthogonal au disque. Le cercle tracé sur les cartes de densité surfacique permet de visualiser le rayon de demi-masse des particules suivies. Les simulations MIRAGE et P sont respectivement montrées dans la colonne de gauche et de droite. La carte de densité en haut à gauche illustre l'identification des particules liées au bulbe stellaire. (Figure du haut) à $t=200$ Myr pour la simulation G1 (à gauche) et la simulation G1P (à droite). (Figure du milieu) à $t=200$ Myr pour la simulation G2 (à gauche) et la simulation G2P (à droite). (Figure du bas) juste après le temps de coalescence t_c pour la simulation G2_G2_90_90_180 (à gauche, $t_c=360$ Myr) et la simulation G2_G2_90_90_180P (à droite, $t_c=434$ Myr).

particules au cours du temps. Dans la figure 4.35, je trace l'évolution de ces profils de densité radiaux pour les simulations G1, G2, et G2_G2_90_90_180 ; simulations pour lesquelles je dispose de leurs homologues de l'échantillon de simulations P et dont je trace aussi le profil ρ_{id} . Pour les galaxies isolées (les quatre figures du haut), le profil de densité initial est choisi au temps $t_i=200$ Myr (courbes noires), l'évolution de ce profil au cours du temps est suivi jusqu'à $t_f=800$ Myr (courbes de couleur). Pour les galaxies en cours de fusion (les deux figures du bas), l'évolution de la distribution des particules stellaires du bulbe est suivi sur une période plus courte, à partir du temps de coalescence des disques t_c jusqu'à 800 Myr. Les profils de densité initial (courbes noires) correspondent donc respectivement aux temps $t_i = t_c=360$ Myr pour la simulation MIRAGE (colonne de gauche) et $t_i = t_c=434$ Myr pour la simulation P, la coalescence étant plus rapide dans les simulations MIRAGE que dans les simulations P parce que la fraction de gaz y est plus importante. On constate que les particules initialement liées au bulbe stellaire dans les simulations MIRAGE (colonne de gauche) sont expulsées des régions centrales vers les régions externes du disque. Cette expulsion est le fait des fusions de granules, qui sont capables de transférer un moment angulaire important aux particules initialement dans le bulbe. À l'inverse, les simulations P (colonne de droite) ne sont pas en mesure de réguler la masse du bulbe car aucune fusion de granule avec éjections n'est observée. Dans les simulations P, la structure du bulbe est plus concentrée et stable au cours du temps.

Dans les figures 4.36, je montre la densité surfacique Σ_{id} des particules stellaires du bulbe identifiées, en utilisant une projection du disque vu par la tranche (plan xz). L'information à deux dimensions contenue dans ces figures permet de situer les zones de migration des particules éjectées. Dans le cas des simulations MIRAGE (colonne de gauche), les particules du bulbe migrent grâce aux fusions de granules riches en gaz vers des trajectoires plus excentriques, et contribuent à la formation d'un ovoïde peuplé d'étoiles vieilles. Cet ovoïde demeure cependant relativement aplati, soulignant de fait un transfert des particules du bulbe vers le disque. La simulation G1 est très efficace pour redistribuer ses étoiles du bulbe vers cette composante plus diffuse du fait des nombreuses fusions de granules très riches en gaz. Les simulations G1P et G2P quant à elles, retiennent la majorité des étoiles du bulbe dans le kpc central, car la formation stellaire non-bridée convertit une grande fraction de gaz en étoiles qui creusent le potentiel gravitationnel au centre, et empêchent l'éjection de baryons malgré un modèle de rétro-action stellaire injectant beaucoup d'énergie thermique.

Dans le cas de la simulation de fusion G2_G2_90_90_180 et G2_G2_90_90_180P, l'identification des particules du bulbe est réalisée sur le pas temps disponible se situant juste après la coalescence des bulbes des deux progéniteurs, respectivement 360 Myr et 434 Myr après les conditions initiales. La simulation G2_G2_90_90_180 est incapable de retenir les étoiles du bulbe, qui subissent plusieurs fusions successives de granules riches en gaz. À la différence des simulations isolées, la simulation de fusion éjecte les étoiles du bulbe en dehors du plan du disque en train de se reformer. En effet, les granules qui fusionnent au centre de masse du système rémanent ont été formés dans deux disques avec un moment angulaire non-colinéaire et éjectent donc le gaz et les étoiles de leurs enveloppes en dehors du plan du disque. De manière opposée, la simulation de fusion G2_G2_90_90_180P concentre la masse dans le centre, et retient toutes les particules dans 1.5 kpc. Cette simulation ressemble plus au schéma canonique de la simulation de fusion qui concentre les baryons au centre de masse pour créer un système sphéroïdal peu étendu. Ce transfert de masse stellaire du bulbe vers le disque est illustré dans la figure 4.37 où l'on constate que les simulations MIRAGE agrandissent systématiquement leurs disques de manière continue, tandis que les simulations P montrent une croissance nulle voir négative.

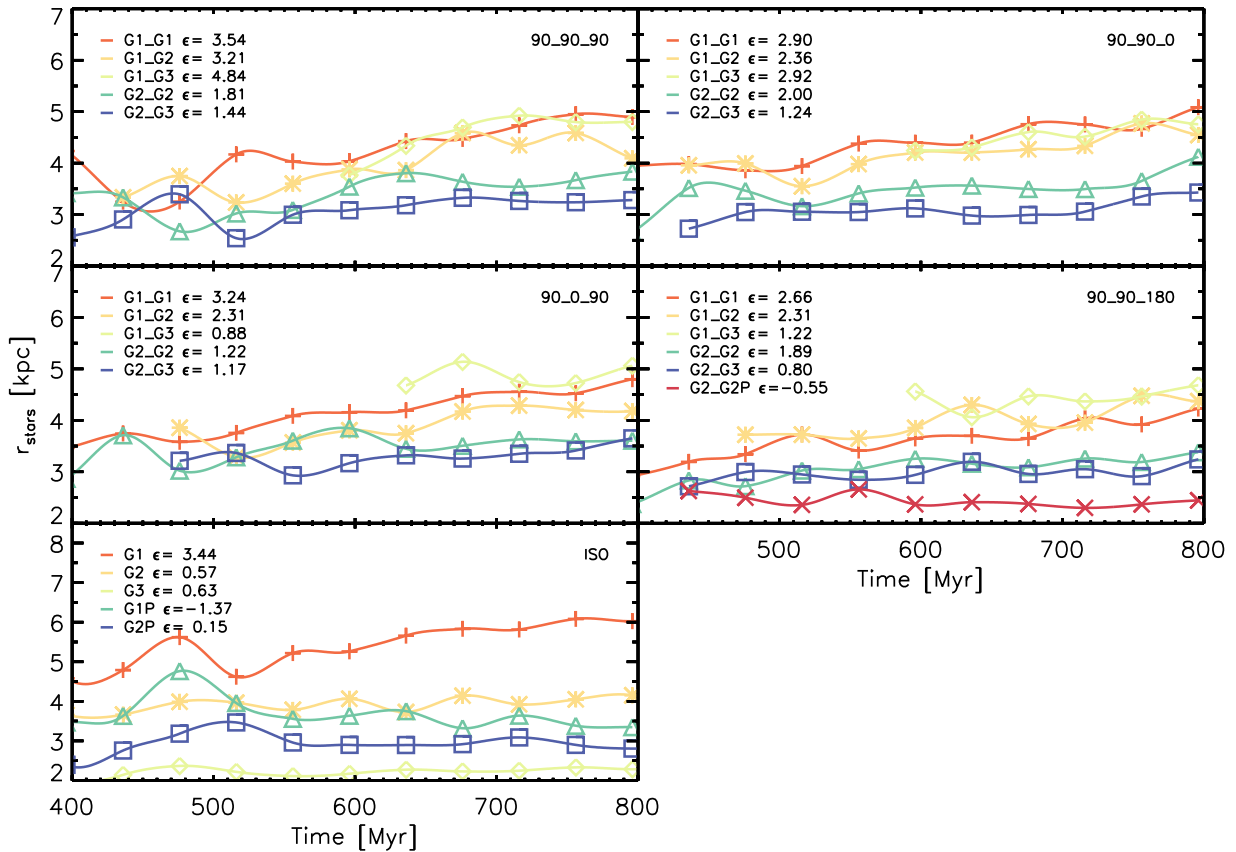


Figure 4.37 – Évolution de l'échelle de longueur dans les simulations MIRAGE et les simulations complémentaires P. La taille des disques stellaires est mesurée grâce à la technique décrite dans la section 3.5. Le taux de croissance des disques stellaires ϵ exprimé en kpc.Gyr^{-1} est mesuré par une régression linéaire sur chacune des courbes. Chaque panneau explore les configurations de masse des progéniteurs indiquée par G_i-G_j pour une orientation initiale des disques fixe indiquée par la concaténation des angles d'orientation de la paire initiale $\theta_1-\theta_2-\kappa$.

4.4.6 Aplatissement du profil de densité du halo

Dans la section précédente nous avons vu que la présence de granules avec une fraction de gaz élevée ($f_g > 80\%$) entretenue par une accrétion de gaz chaud du milieu circumgalactique permet de redistribuer les étoiles du bulbe dans le disque et l'ovoïde stellaire grâce au mécanisme de fusion et aux forces de marées attenantes. Ce mécanisme, étendu à la matière non-baryonique et non collisionnelle constitué par la matière noire, permet notamment d'aborder le problème des galaxies produites par les simulations cosmologiques qui présentent des courbes de rotations trop piquées dans les régions centrales du fait d'un potentiel gravitationnel très profond au centre du disque, contrairement aux observations qui montrent des courbes de rotation dont la croissance dans les parties internes est modérée, indiquant que le profil de densité est plutôt uniforme que cuspidé dans ces régions centrales ($\sim 1/3$ à $1/2 r_d$, r_d étant l'échelle de longueur du disque).

La fusion des granules au centre de masse des galaxies est aussi un processus à même de chauffer les parties centrales du halo de matière sombre, de manière strictement identique aux étoiles. Les particules de matière sombre présentes dans le kilo-parsec central sont déviées sur des orbites plus excentriques, ce qui contribue à abaisser et aplatir le profil central de la densité volumique de matière noire. Inoue et Saitoh (2011) remarquent la présence d'une "érosion" de la pente du profil de densité du halo dans des disques soumis à une phase d'instabilité gravitationnelle, i.e. lorsque les galaxies sont dans leur phase "granulaires". La coalescence des granules chauffe et lisse par friction dynamique le pic de densité de matière noire. De surcroît au chauffage par friction dynamique, les violentes oscillations dans le potentiel gravitationnel dues à l'éjection rapide de grandes quantités de gaz sont aussi en mesure de chauffer gravitationnellement le noyau du halo par impulsions. L'échange violent de moment angulaire entre le coeur des granules et leur enveloppe de gaz éjectée à des vitesses angulaires très élevées ($> 300 \text{ km.s}^{-1}$) reproduit le schéma de chauffage par impulsions. Un processus similaire est modélisé numériquement dans les galaxies naines qui font alors l'expérience d'une rétro-action stellaire (supernova, pression radiative, etc.) injectant de l'énergie de manière cyclique de façon à ce que la structure du halo s'en retrouve affectée (Governato et al., 2010; Teysier et al., 2013). La figure 4.38 présente l'évolution des profils de densité des halos de matière sombre. On constate dans le cas des simulations de l'échantillon MIRAGE que les profils s'abaissent et s'aplatissent, contrairement à ceux des simulations P qui sont tous dynamiquement stabilisés. Dans le cas de la fusion, les profils sont mesurés seulement après la coalescence des bulbes des deux progéniteurs. Même si il est difficile de distinguer l'origine exacte de cette érosion, la combinaison de des deux processus (friction dynamique et chauffage impulsif) semble être à l'œuvre dans les simulations de l'échantillon MIRAGE. La migration des granules vers le bulbe central chauffent par friction dynamique le halo. Lorsque les granules riches en gaz fusionnent, les forces de marées produisent une forte impulsion dans le potentiel gravitationnel, ce qui ajoute un chauffage gravitationnel supplémentaire aux particules de matière sombre situées dans les régions centrales. Inoue et Saitoh (2011) constatent au demeurant que la formation d'un bulbe stellaire massif peut inverser le mécanisme et reformer un profil de densité piqué. Les mécanismes de formation du bulbe stellaire et du noyau de matière sombre sont intimement liés. L'aptitude du mécanisme de fusion des granules riches en gaz à contrôler à la fois la croissance du bulbe et le profil de matière sombre permet d'apporter de nouveaux éléments au problème "cusp/core" pour les galaxies très massives, qui restent une épine dans le pied des modélisations cosmologiques numériques des galaxies.

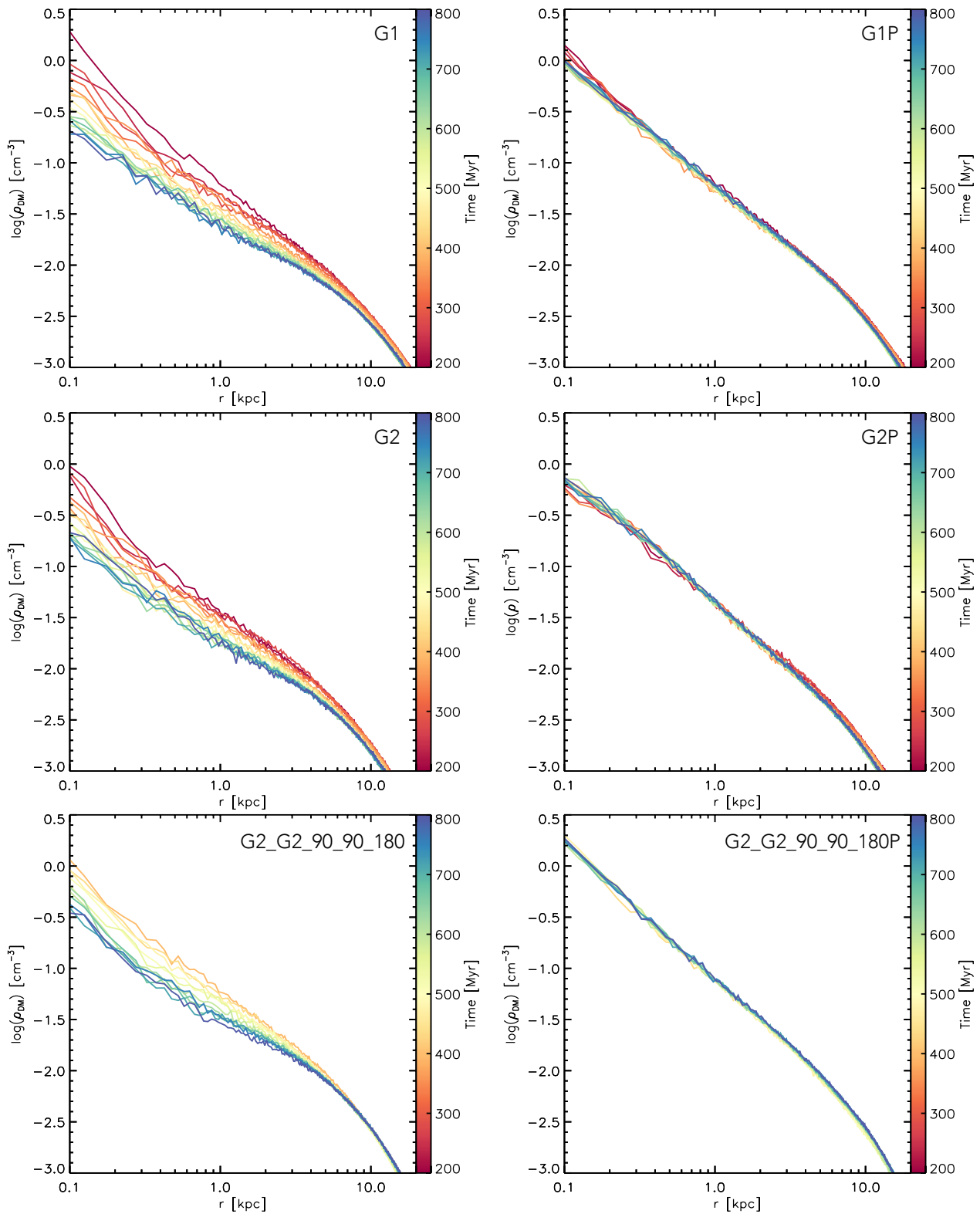


Figure 4.38 – Évolution des profils de densité des halos de matière noire. De gauche à droite : en haut, simulations G2_G2_90_90_180 et G2_G2_90_90_180P, en bas, simulations G1 et G1P. Le pas de temps est représenté par la couleur de la courbe de densité. Les simulations P montrent des profils stabilisés tout au long des 800 Myr simulés, tandis que les simulations MIRAGE montrent une érosion du profil interne et une redistribution vers les régions externes de la masse.

Pseudo-observations de l'échantillon



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Ce chapitre présente une analyse comparative de l'échantillon d'observations spectroscopiques MASSIV et de l'échantillon de simulations numériques MIRAGE, tous deux introduits dans les chapitres précédents. Une comparaison qualitative et quantitative des deux échantillons nécessite de passer par une étape de création de pseudo-observations¹ à partir des données brutes de simulations. Ces simulations d'observations sont ensuite analysées et réduites en utilisant les outils développés pour les observations. Ce chapitre a pour vocation de faire le lien entre tous les travaux engagés durant cette thèse, en mêlant l'analyse des observations à celle des simulations.

5.1 Méthode

Une des grandes difficultés concernant l'analyse des simulations numériques est de réussir à replacer les données numériques dans un contexte observationnel. Dans ce but, j'ai développé un outil permettant la création de pseudo-observations. Ces pseudo-observations permettent de comparer des données observationnelles à des simulations numériques (avec toutes les précautions d'usage à adopter dans ce type d'analyse) pour des échantillonnages et des résolutions spatiales et spectrales comparables. Dans cette section, je présente l'outil

1. "mock observations" en anglais.

Aurora qui est le fruit de ce travail. Ce programme est capable de lire des fichiers de simulations de type GADGET2 (Springel, 2005) et de produire différentes cartes d'observables à partir de quantités physiques. En utilisant la bibliothèque de conversion de formats UNSIO², il est possible de traiter de nombreux autres types de formats de simulations avec le programme Aurora. Dans le cadre de cette thèse, j'ai travaillé principalement sur des formats de type RAMSES, convertis au format GADGET2 grâce à cette bibliothèque. Il est à noter que le code Aurora fonctionne à la fois sur des simulations utilisant une implémentation Lagrangienne (e.g. SPH) et Eulerienne (e.g. AMR) du solveur hydrodynamique. Le programme Aurora se subdivise en trois modules distincts que je décris dans les paragraphes suivants.

5.1.1 Cartes de quantités physiques

Afin de pouvoir comprendre le contenu d'un fichier de sortie d'une simulation numérique, il est nécessaire de disposer d'un outil permettant d'afficher différentes quantités physiques sur des cartes 2D. Un module du programme Aurora permet de créer de telles cartes. Ce programme utilise toujours en entrée des données transcrites dans le format des particules Lagrangiennes, même si le code d'origine ayant produit la simulation utilise un schéma Eulerien. De manière générale, une quantité hydrodynamique est projetée sur une carte 2D à une résolution donnée en : (i) orientant la simulation suivant la géométrie désirée, (ii) en calculant les coordonnées de chaque particule sur la matrice de la carte, (iii) en calculant au choix la valeur moyenne/médiane/maximale/minimale de la quantité hydrodynamique dans chaque pixel de la carte.

Lorsque l'échantillonnage de la carte que l'on souhaite créer se rapproche de la résolution de la simulation numérique, un élément hydrodynamique peut alors contribuer à plusieurs pixels de la carte. Déterminer cette contribution pour chaque élément dans l'espace mathématique des nombres réels peut se révéler extrêmement coûteux, c'est pour cela qu'il est intéressant de passer dans l'espace de Fourier pour traiter ce problème. En effet, à chaque élément hydrodynamique d'une simulation est associée une longueur qui quantifie le volume dans lequel sont estimées les quantités physiques (densité, température, vitesse, etc.) : c'est la résolution spatiale locale. En traitant ces éléments hydrodynamiques par familles de volumes, on peut estimer pour une famille de volume i la contribution spatiale en convoluant la carte de la quantité physique considérée par un noyau Gaussien dont la largeur à mi-hauteur est proportionnel la résolution spatiale locale $\epsilon(i)$. Un facteur trois permet d'étaler spatialement l'information hydrodynamique de manière optimale, c'est-à-dire avec une dégradation minimale de la résolution de la projection. Cette technique rapide permet d'estimer dans quels pixels de la carte considérée contribue un élément hydrodynamique. En sommant toutes les sous-cartes convoluées, on obtient notre carte finale :

$$\mathbb{M}_{\text{corr}}(x, y) = \sum_i \left[\mathbb{M}(x, y, i) \otimes_{xy} \mathbb{G}(x, y, \epsilon(i)) \right], \quad (5.1)$$

avec $\mathbb{M}_{\text{corr}}(x, y)$ la carte 2D finale corrigée du faible échantillonnage de la simulation dans certains volumes, $\mathbb{M}(x, y, i)$ la carte 2D contenant seulement des particules de gaz modélisant un volume caractérisé par $\epsilon(i)$, et $\mathbb{G}(x, y, \epsilon(i))$ le noyau Gaussien de convolution. La conservation des quantités physiques est assurée un noyau Gaussien normalisé. De cette manière, on interpole de manière efficace les quantités physiques dans les espaces "vides" entre les po-

2. La bibliothèque UNSIO est disponible à l'adresse <http://projets.lam.fr/projects/unsio>

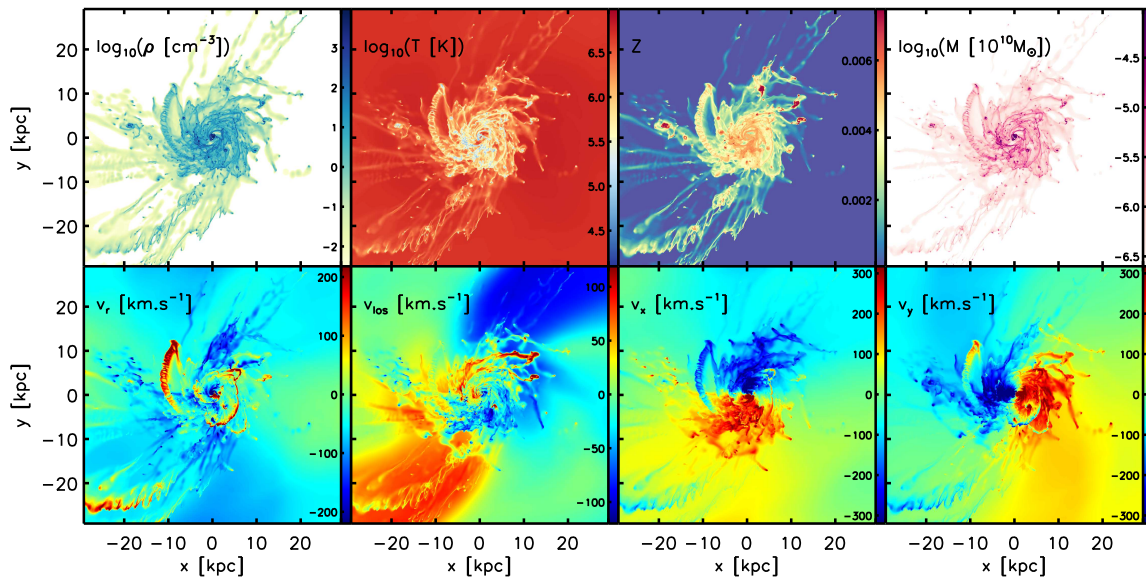


Figure 5.1 – Exemple de cartes de différentes quantités physiques hydrodynamiques associées au gaz et produites par le programme Aurora. Les cartes présentées sont calculées pour le modèle G1_G1_90_90_180 de l'échantillon MIRAGE, à $t=680$ Myr pour une projection de face du disque reformé après une récente fusion. Sur la rangée du haut, on trouve de gauche à droite : une carte de densité, une carte de température, une carte de métallicité et une carte de masse. Sur la rangée du bas, on trouve de gauche à droite : une carte des vitesses radiales cylindriques dans le plan de projection qui est coplanaire au plan du disque dans ce cas, une carte des vitesses le long de la ligne de visée perpendiculaire au disque, une carte de vitesses de la composante x du plan de projection, une carte de vitesses de la composante y du plan de projection.

sitions des centres des particules de gaz. Un exemple de cartes dérivées par le programme Aurora est présenté dans la figure 5.1.

5.1.2 Pseudo-observations photométriques

Le second module du programme Aurora permet de créer des observations photométriques. Pour ce faire, il est nécessaire de fournir au module un modèle d'émission du continu stellaire ainsi qu'une transmission pour le filtre photométrique considéré. Dans le cadre de ce travail, j'utilise essentiellement le modèle STARBURST99³ (Leitherer et al., 1999) qui convient à ce type d'exercice, mais n'importe quel modèle de SED est utilisable du moment qu'il se présente sous le bon format de fichier. La figure 5.2 présente un exemple de modèle de SED utilisé dans ce travail pour une population stellaire de $10^6 M_{\odot}$ avec une fonction de masse initiale de Salpeter (1955).

En fournissant un fichier contenant la transmission du filtre photométrique souhaité pour produire la pseudo-observation, il est possible d'intégrer le flux lumineux d'une population stellaire donnée. Ce module du programme Aurora récupère donc la liste de particules stellaires contenues dans le fichier de simulation. Chaque particule stellaire est assimilée à une

3. <http://www.stsci.edu/science/starburst99/docs/default.htm>

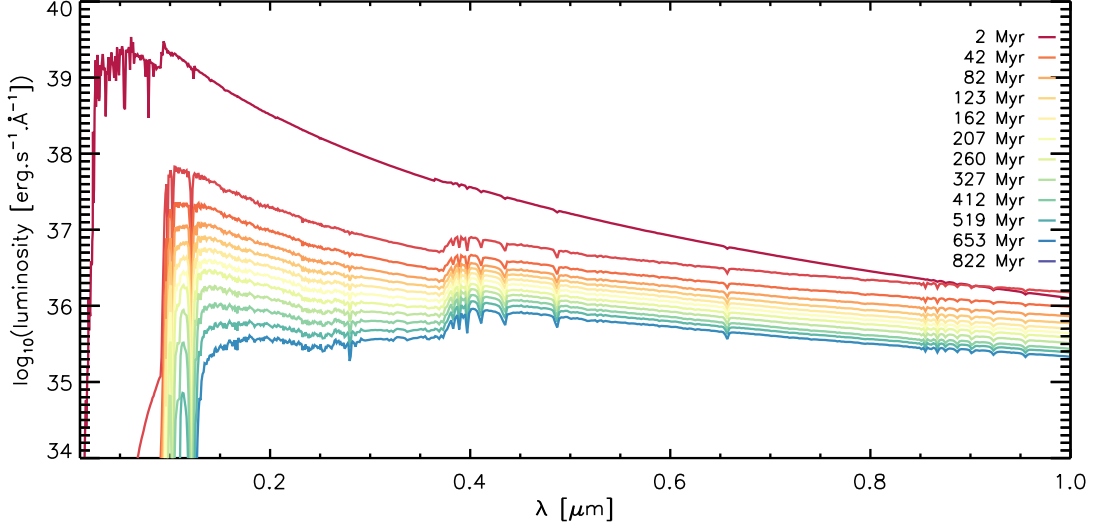


Figure 5.2 – Distribution d'énergie spectrale pour une population stellaire de $10^6 M_{\odot}$ avec une IMF de Salpeter (1955) ($\alpha = -2.35$) et une métallicité $Z = 0.04$ calculée par le modèle STARBURST99. La couleur encode l'âge de la population stellaire modélisée par la courbe de distribution spectrale d'énergie.

population d'étoiles de différentes masses dont le spectre d'émission dans le référentiel au repos peut être simulé par un modèle de SED $\ell_{\star,0}$. Ce spectre dépend de l'âge t de formation de cette population stellaire. On applique une extinction au spectre du continu stellaire en utilisant la courbe d'obscurcissement $k^e(\lambda)$ de Calzetti et al. (2000) :

$$\ell_{\star,0,ext}(\lambda, t) = \ell_{\star,0}(\lambda, t) \times 10^{0.4E(B-V)k(\lambda)}, \quad (5.2)$$

avec $E(B-V)$ le paramètre d'extinction de la poussière interstellaire, librement spécifié dans les paramètres d'entrée du module. Pour une galaxie située à un décalage spectral z , la densité spectrale à une longueur d'onde observée λ' s'écrit :

$$\ell_{\star}(\lambda', t, z) = \frac{\ell_{\star,0,ext}(\lambda, t)}{1+z}, \quad (5.3)$$

avec $\lambda' = \lambda(1+z)$ décalée pour tenir compte de l'expansion de l'univers entre le décalage spectral z et l'époque actuelle $z = 0$. Le modèle de SED ayant l'âge t le plus proche est utilisé pour estimer $\ell_{\star,0}(\lambda, t)$. La luminosité intégrée dans un filtre photométrique X est obtenue en réalisant l'intégration numérique :

$$L_{\star,X}(t, z) = \int_{\mathbb{T}_{X,min}}^{\mathbb{T}_{X,max}} \ell_{\star}(\lambda, t, z) \mathbb{T}_X(\lambda) d\lambda, \quad (5.4)$$

avec $\mathbb{T}_X(\lambda)$ la transmission du filtre à la longueur d'onde λ , $\mathbb{T}_{X,min}$ et $\mathbb{T}_{X,max}$ les limites inférieure et supérieure en longueur d'onde du filtre X . Finalement, le flux intégré du continu stellaire dans le filtre X est obtenu grâce à l'équation :

$$F_{\star,X} = \frac{L_{\star,X}}{4\pi d_l^2} \quad (5.5)$$



Figure 5.3 – Pseudo-observations photométriques de la simulation G2_G2_90_90_180P réalisées aux pas de temps avoisinant la coalescence. Les images composites présentées sont réalisées à partir de trois pseudo-observations qui simulent l’instrument SDSS dans les filtres ugr en projetant la simulation à $z=0.01$, et en utilisant un modèle de SED calculé avec STARBURST99.

avec d_l la distance lumineuse cosmologique, calculée pour des paramètres cosmologiques de Planck-2013 ($H_0 = 67.8 \text{ km.s}^{-1}.\text{Mpc}^{-1}$, $\Omega_m=0.307$, et $\Omega_\Lambda=0.691$). Le calcul des distances lumineuses cosmologiques est assuré par la bibliothèque python `astrophy`⁴. Une cosmologie différente (e.g. WMAP-9) peut être spécifiée dans les paramètres d’entrée du programme.

Afin de créer une pseudo-observation photométrique, il est nécessaire de spécifier la résolution et l’échantillonnage spatial de la carte simulée. De même que pour construire les cartes de quantités hydrodynamiques, on oriente dans un premier temps la simulation selon la ligne de visée désirée puis on détermine les coordonnées des particules stellaires sur la carte. La valeur en chaque pixel est le résultat de la somme des contributions en flux des particules stellaires localisées dans ce pixel. On applique un lissage avec un noyau Gaussien pour simuler la PSF instrumentale.

Enfin, il est possible d’injecter dans la pseudo-observation photométrique un bruit Gaussien dont le niveau est calculé grâce à la magnitude AB (Oke, 1974) limite, couramment utilisée dans les grand relevés photométriques. La dispersion du bruit Gaussien exprimée en unité de flux par pixel s’écrit :

$$\sigma_{\star,X} = 10^{\left(\frac{-48.6 - X_{AB,limit}}{2.5}\right)} \Delta x^2 \int_0^{+\infty} \left(-\frac{\mathbb{T}_X(\lambda)c}{\lambda^2} \right) d\lambda \quad (5.6)$$

avec $X_{AB,limit} = -2.5 \log_{10} f_{\nu,limit} - 48.60$ la magnitude AB limite de détection pour une densité spectrale de flux limite $f_{\nu,limit}$ et Δx la taille du pixel du pseudo-observation. L’intégrale de la membre de droite de l’équation correspond à l’intégration de la transmission du filtre X sur le domaine de fréquence $\int_0^{+\infty} \mathbb{T}_X(\nu) d\nu$.

5.1.3 Pseudo-observations spectroscopiques

Le dernier module du programme Aurora permet de simuler des observations spectroscopiques 3D, i.e. de générer des pseudo-cubes spectroscopiques. Ces cubes contiennent

4. <http://www.astrophy.org>

deux dimensions spatiales et une dimension spectrale qui permet de dériver des informations spectroscopiques dans différentes régions spatiales de l'objet (e.g. position de raies, largeurs de raies, rapports de raies, etc.). Dans le cadre de cette thèse, je me suis essentiellement concentré sur le cas des observations de la raie d'émission $H\alpha$ qui est la raie principalement utilisée dans l'échantillon MASSIV, mais de nombreuses autres raies d'émission (e.g. [OII], [OIII], [NII], etc.) ou d'absorption (e.g. $H\beta$, Mg b , Fe) sont utilisées dans ce type d'observations. Afin de construire un pseudo-cube, je calcule pour chaque élément hydrodynamique d'une simulation donnée une luminosité $H\alpha$ en suivant la description de Rozas et al. (2006) pour les régions HII. Cette luminosité dépend de la densité électronique et de la température T de la particule de gaz :

$$L_{H\alpha} = V n_e^2 h \nu_{H\alpha} \alpha_{H\alpha}^{eff}(T, H_0), \quad (5.7)$$

avec V le volume physique de la particule de gaz, h la constante de Planck, $\nu_{H\alpha}$ la fréquence des photons $H\alpha$ au repos, n_e le nombre d'électrons disponibles pour la recombinaison dans le volume modélisé par la particule de gaz et $\alpha_{H\alpha}^{eff}(T, H_0) = 2.59 \times 10^{-13} (T/10^4)^{-0.845}$ le taux de recombinaison effectif de l'atome d'hydrogène pour la raie d'émission Balmer- α pour un volume de gaz à la température T (Osterbrock et Ferland, 2006). Le flux intégré $H\alpha$ d'une particule de gaz est calculé grâce à la distance lumineuse cosmologique d_l :

$$F_{H\alpha} = \frac{L_{H\alpha}}{4\pi d_l^2}. \quad (5.8)$$

Le flux $H\alpha$ est distribué sur le domaine spectral en utilisant une fonction Gaussienne :

$$f_{H\alpha}(\lambda) = \frac{F_{H\alpha}}{\sigma_\lambda \sqrt{2\pi}} \exp\left(-\frac{(\lambda - \lambda_{H\alpha})^2}{\sigma_\lambda^2}\right), \quad (5.9)$$

avec $\lambda_{H\alpha}$ la longueur d'onde de la raie $H\alpha$ décalée spectralement par effet Doppler (distance cosmologique et mouvements propres) et σ_λ la largeur spectrale de la raie d'émission. La largeur de la raie d'émission associée à chaque particule de gaz est calculée en considérant que le support thermique domine à l'intérieur du volume modélisé par la particule de gaz, ce qui implique une dispersion en terme de vitesses radiales :

$$\sigma_v = \frac{\sigma_\lambda \lambda_{H\alpha}}{c} = c \sqrt{\frac{k_B T}{m_H c^2}}, \quad (5.10)$$

avec c la vitesse de la lumière, k_B la constante de Boltzmann, m_H la masse de l'hydrogène et T la température de la cellule. Une température de 10^4 K, commune dans les disques des galaxies spirales locales induit une dispersion des vitesses de 10 km.s^{-1} .

Enfin, l'extinction du flux par la poussière interstellaire est calculée en utilisant la densité intrinsèque de flux "starburst" $f_{H\alpha,ext}(\lambda)$ via :

$$f_{H\alpha,ext}(\lambda) = f_{H\alpha}(\lambda) \times 10^{0.4E(B-V)k_{H\alpha}}, \quad (5.11)$$

avec $k_{H\alpha} = 2.517$ le coefficient d'obscurcissement pour la raie d'émission $H\alpha$ (Boselli, 2011). Cette correction peut être appliquée dans le référentiel de l'observateur car $k_{H\alpha}$ ne dépend pas de la longueur d'onde.

Afin de créer un pseudo-cube, il est nécessaire de spécifier la résolution et l'échantillonnage dans les directions spatiales et spectrales. Les premières étapes sont les mêmes que celle décrites pour les pseudo-observations photométriques (orientation de la simulation et calcul des coordonnées des particules). Contrairement aux pseudo-observations photométriques, chaque particule de gaz modélise un volume physique de la simulation. Dans le cas de pseudo-observations à forte résolution spatiale (de l'ordre de la résolution de la simulation), il est nécessaire d'étendre l'information de flux dans les pixels échantillonnant le volume de la particule de gaz considérée : pour chaque famille de taille de particules de gaz, le spectre en chaque pixel est le résultat de la somme des spectres des particules de gaz localisées dans ce pixel. L'extension de l'information spatiale est effectuée en utilisant la même méthode que celle utilisée pour créer les cartes de quantités hydrodynamiques (cf. équation 5.1). Les résolutions spatiale et spectrale des pseudo-données sont simulées en appliquant un lissage spatial et un lissage spectral avec des noyaux Gaussiens.

Ce module est de plus en mesure de calculer le niveau de l'émission du continu stellaire en procédant de manière similaire au module de pseudo-observations photométriques. Pour chaque spectre du pseudo-cube, la contribution au flux lumineux du continu stellaire est estimée en interpolant les données du modèle de SED. Finalement, un bruit Gaussien peut être injecté dans le pseudo-cube afin de n'observer que le signal détectable en conditions observationnelles. De manière optionnelle, il est possible de remplacer le bruit Gaussien par un bruit Poissonien qui est plus réaliste pour modéliser une réalisation aléatoire de photon. Un exemple de champ de vitesses déduit d'un cube λ modélisant une observation SINFONI à $z=1.7$ grâce à cette méthode est présenté dans la figure 5.4.

5.1.4 Transfert radiatif

Le programme Aurora se veut un outil versatile et rapide, capable de générer de nombreuses pseudo-observations afin de pouvoir réaliser des analyses statistiques. C'est pourquoi Aurora fait l'approximation d'un milieu interstellaire transparent aux rayonnements, ce qui est une approximation du premier ordre. En effet, afin de tenir compte des processus d'absorption, de diffusion et de ré-émission des photons, il devient nécessaire d'introduire du transfert radiatif, or, modéliser numériquement de tels processus est très coûteux.

Le fait d'inclure un transfert radiatif dans la production de pseudo-observations affecte nécessairement le flux, la position, la largeur des raies d'émission ainsi que le niveau de continu stellaire. Les pseudo-observations produites par Aurora sont donc à considérer comme des approximations du premier ordre dont on peut déjà extraire une grande quantité d'informations scientifiquement pertinentes, mais il convient d'être conscient de cette limitation lors de l'analyse. L'ajout d'un module de transfert radiatif simplifié est prévu comme perspective proche pour le programme Aurora afin de quantifier l'écart au cas du milieu optiquement fin.

5.2 Paramétrage des pseudo-observations SINFONI

Afin de pouvoir réaliser une étude comparative entre l'échantillon MASSIV et les simulations de l'échantillon MIRAGE, j'ai produit une série de pseudo-observations avec les modules du programme Aurora introduits précédemment. Le but de cette approche scientifique est de disposer d'une base de données de pseudo-observations imitant des données SINFONI ainsi que des données photométriques CFHT (bande I) afin de pouvoir les compa-

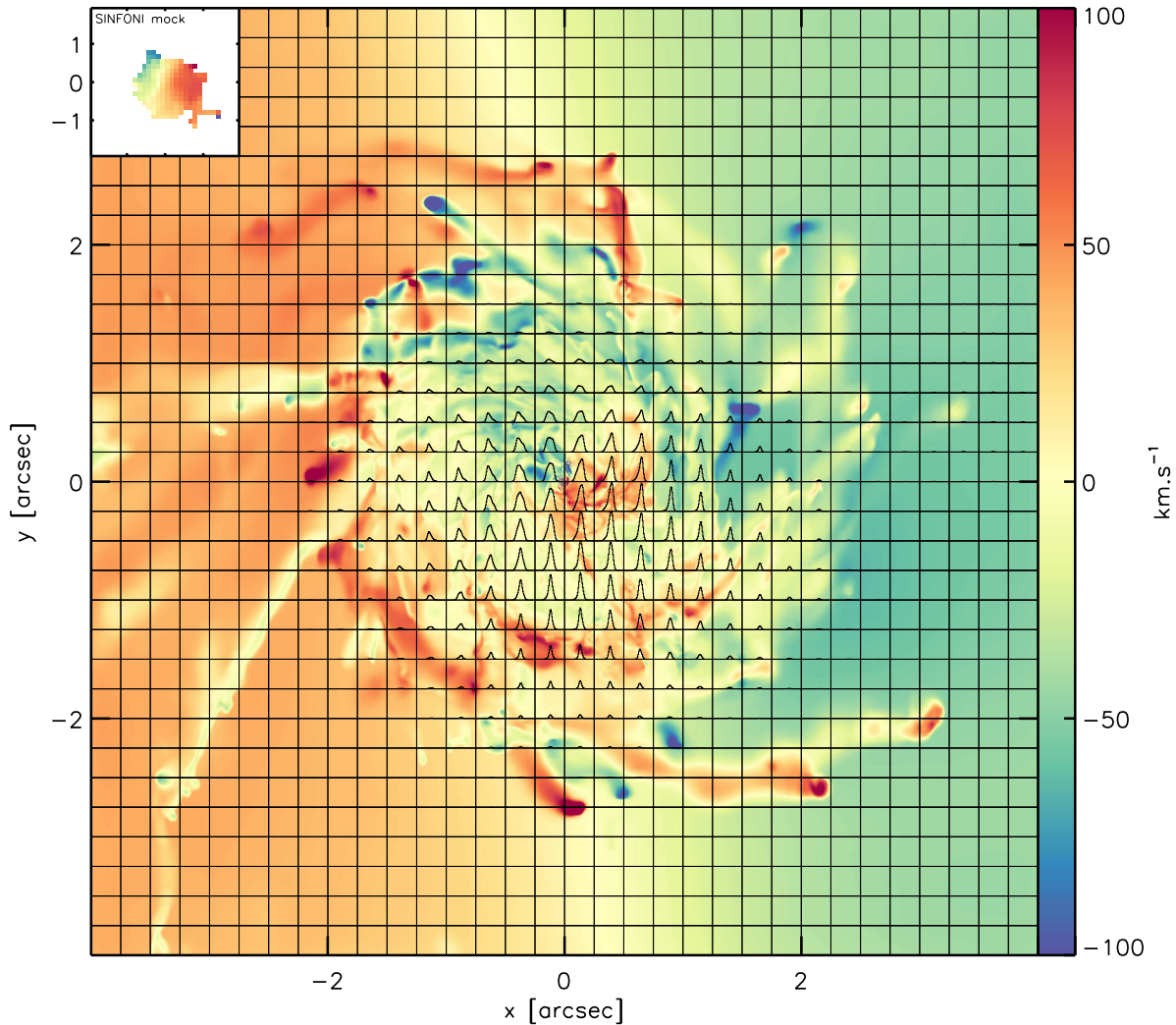


Figure 5.4 – Champ de vitesses de la simulation G2_G2_90_90_90 à $t=760$ Myr. La ligne de visée “los6” est utilisée pour réaliser la projection. La carte de vitesses haute résolution est réalisée avec le programme Aurora qui calcule pour chaque élément spatial une vitesse moyenne pondérée par la masse des cellules AMR du pixel. Par dessus cette carte, on affiche les raies d’émission $H\alpha$ normalisées du pseudo-cube spectroscopique SINFONI projeté à $z=1.7$. Chaque raie affichée couvre 0.125 secondes d’arc, soit la taille du pixel SINFONI. La taille du champ couvert par le pseudo-cube et la carte de vitesse haute résolution est identique (4 secondes d’arc pour $z=1.7$). Le champ de vitesses du pseudo-cube retrouvé par les outils d’analyses MASSIV est montré dans la vignette en haut à gauche.

rer aux données spectroscopiques et morphologiques des galaxies à $z > 1$ observées dans le cadre de MASSIV. Bien que ce chapitre se concentre sur une comparaison aux données MASSIV introduites dans le chapitre 1, cette méthode a pour vocation d’être utilisée pour n’importe quel échantillon spectroscopique de galaxies à grand décalage spectral.

En premier lieu, la géométrie des projections est définie grâce à un jeu de 11 lignes de visées distribuées de manière régulière sur la moitié de la surface d’une sphère centrée sur la galaxie (figure 5.5). En effet, en faisant l’hypothèse d’un milieu interstellaire transparent

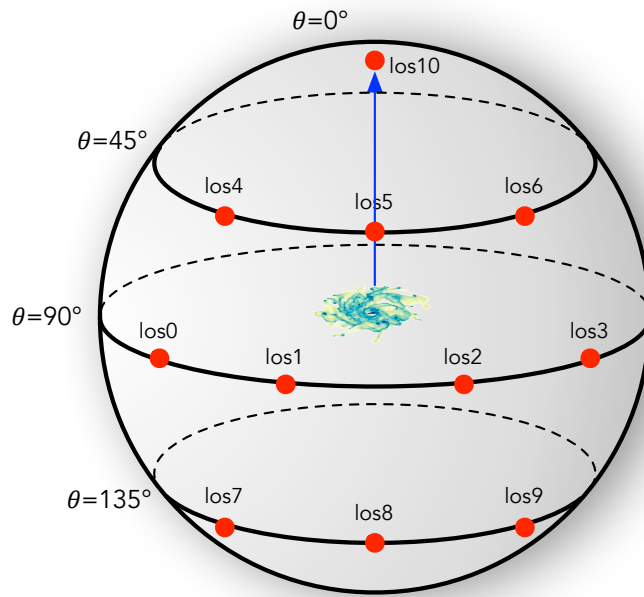


Figure 5.5 – Géométrie des angles de projections choisis pour construire les pseudo-observations SINFONI. Chaque point rouge représente une orientation pour une ligne de visée. Entre chacune de ces lignes de visées, on impose un angle solide minimal de $\Omega_{min} \sim 0.46$ stéradian. Comme le milieu interstellaire est supposé transparent aux rayonnements, il n'est pas nécessaire de couvrir l'intégralité de la surface de la sphère du fait des symétries centrales qui existent. Le vecteur de moment angulaire du disque (vecteur bleu) est aligné avec la ligne de visée "los10".

aux rayonnements, les symétries justifient le fait de n'échantillonner qu'une demi-sphère. Par ailleurs, le fait d'échantillonner la demi-sphère de manière régulière permet d'éviter des lignes de visée trop proches pour lesquelles on s'attendrait à avoir des pseudo-observations similaires. Ainsi, chacune des 11 lignes de visées est séparée des lignes de visée voisines avec angle solide minimal de $\Omega_{min} \sim 0.46$ stéradian. Les angles définissant les lignes de visées sont présentés dans la Table 5.1.

	los0	los1	los2	los3	los4	los5	los6	los7	los8	los9	los10
ϕ (longitudes)	22.5	67.5	112.5	157.5	26.4	90	153.5	26.4	90	153.5	0
θ (latitudes)	90	90	90	90	45	45	45	135	135	135	0

Table 5.1 – Angles en degrés définissant les 11 lignes de visée utilisées pour créer la série de pseudo-observations SINFONI à partir des simulations numériques de l'échantillon MIRAGE.

J'ai choisi de produire 11 pseudo-observations SINFONI tous les 40 Myr entre 200 et 800 Myr pour chacune des simulations de l'échantillon MIRAGE. Avec un total de 23 simulations dans l'échantillon MIRAGE, ce paramétrage permet de produire un total de 4048 pseudo-observations SINFONI et CFHT. Avant chaque projection, je me suis assuré d'avoir replacé le centre du disque tel que défini dans le chapitre 3 à la position cartésienne (0,0,0). De plus,

j'ai aligné le vecteur de moment angulaire des étoiles jeunes avec l'axe z du système de coordonnées cartésiennes afin d'être assuré d'avoir un disque vu de face avec l'angle de projection $\theta = 0^\circ$ et un disque vu par la tranche avec l'angle de projection $\theta = 90^\circ$. Dans le cas des simulations de fusion où la coalescence des deux disques n'est pas complète, j'ai utilisé la correction pour l'alignement du vecteur de moment angulaire avec l'axe z du cas le plus proche temporellement de la coalescence. De ce fait, les pseudo-observations des simulations de fusion pre-coalescence ne montrent pas nécessairement des disques vus de face pour $\theta = 0^\circ$. Les différents paramètres de projection pour les pseudo-observations SINFONI et CFHT sont présentés dans la Table 5.2.

	SINFONI	CFHT I-band
1. z		1.7
1. d_l [Mpc]		12994
1. scale [kpc/"]		8.64
2. Δx ["]	0.125	0.205
3. $\Delta \lambda$ [Å]	1.95	-
4. S ["]	0.60	0.80
5. R	2500	-
6. $n_{spatial}$	32×32	30×30
7. $n_{spectral}$	48	-
8. R_V	-	4.04
9. $E(B-V)$		0.10
10. ρ_{min} [cm ⁻³]	0.1	-
11. ρ_{max} [cm ⁻³]	100	-
12. $m_{AB,limit}$	-	25.0
13. σ_{noise} [erg.s ⁻¹ .cm ⁻² .μm ⁻¹]	1.0×10 ⁻¹⁵	-

Table 5.2 – Paramètres de projection pour les pseudo-observations SINFONI et CFHT. 1. Décalage spectral, distance lumineuse et facteur d'échelle. 2. Échantillonnage spatial. 3. Échantillonnage spectral. 4. Résolution spatiale (largeur à mi-hauteur). 5. Résolution spectrale. 6. Nombre de pixels spatiaux. 7. Nombre de canaux spectraux. 8. Coefficient d'absorption. 9. Extinction. 10. Densité minimale du gaz participant à l'émission $H\alpha$. 11. Densité maximale du gaz participant à l'émission $H\alpha$. 12. Magnitude AB limite. 13. Limite de detection spectrale.

Les cartes des champs de vitesses, des champs de dispersion de vitesses, et les cartes de flux stellaire dans le filtre I des pseudo-observations produites pour la simulation G1_G1_90_90_0 sont présentés respectivement dans les figures 5.6 et 5.7 et 5.8. La totalité des champs de vitesses des pseudo-observations SINFONI est présentée dans l'annexe A de ce manuscrit.

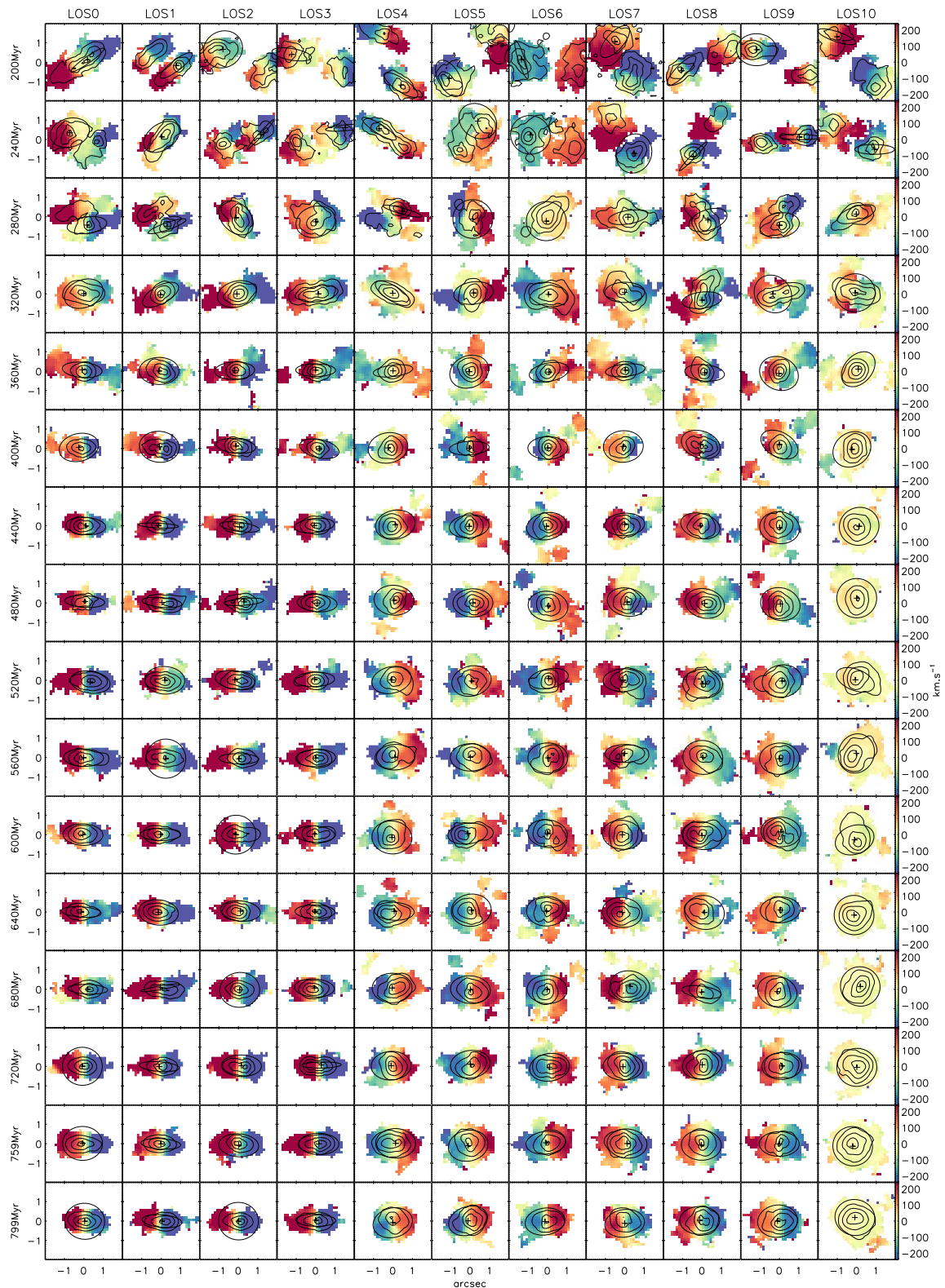


Figure 5.6 – Champs de vitesses des pseudo-observations SINFONI de la simulation G1_G1_90_90_0 de l'échantillon MIRAGE. Chaque ligne présente un temps spécifique dans la simulation, chaque colonne présente une ligne de visée différente pour la pseudo-observation. Sur chaque champ de vitesses, les contours issus de la carte de flux H α sont tracés. L'ellipse centrée sur une croix présente le résultat de l'ajustement des paramètres géométriques grâce à la routine KINEGEN présentée dans le chapitre 1.

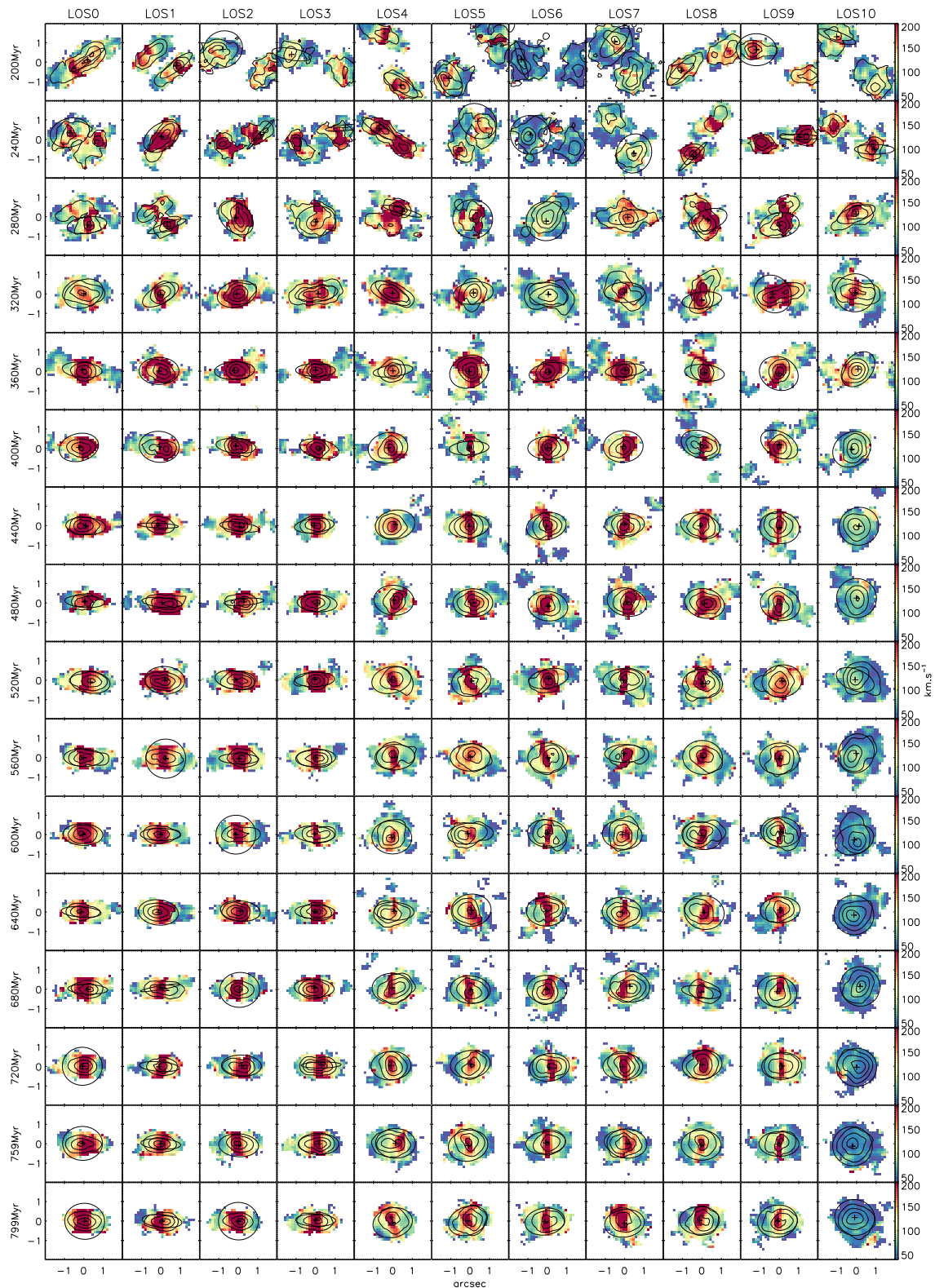


Figure 5.7 – Champs de dispersion de vitesses des pseudo-observations SINFONI de la simulation G1_G1_90_90_0 de l'échantillon MIRAGE. Chaque ligne présente un temps spécifique dans la simulation, chaque colonne présente une ligne de visée différente pour la pseudo-observation. Sur chaque champ de vitesses, les contours issus de la carte de flux H α sont tracés. L'ellipse centrée sur une croix présente le résultat de l'ajustement des paramètres géométriques grâce à la routine KINEGEN.

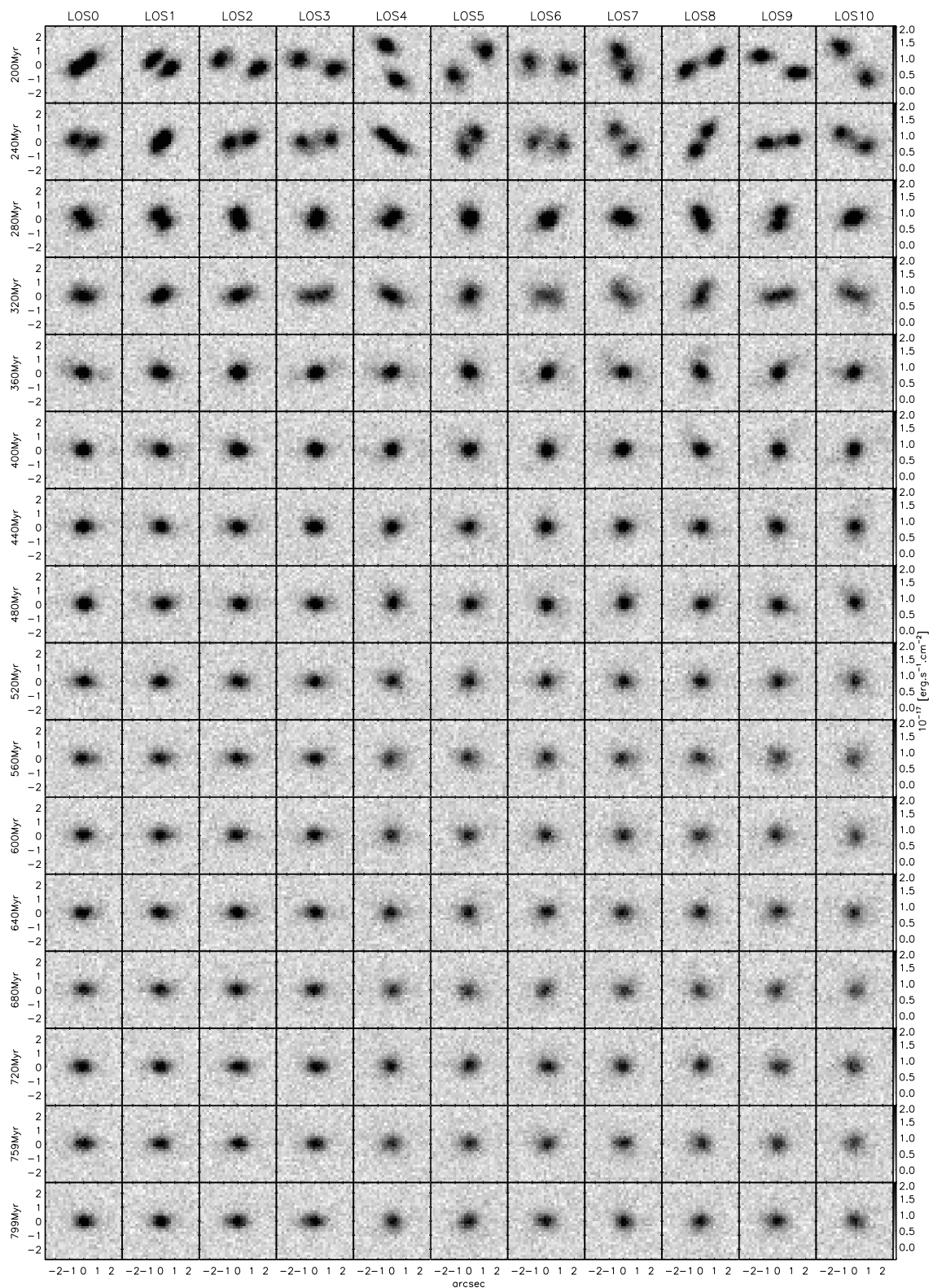


Figure 5.8 – Carte de flux stellaire des pseudo-observations CFHT en bande I de la simulation G1_G1_90_90_0 de l'échantillon MIRAGE. Chaque ligne présente un temps spécifique dans la simulation, chaque colonne présente une ligne de visée différente pour la pseudo-observation.

5.3 Analyse cinématique

5.3.1 Ajustement des paramètres géométriques

Chacun des pseudo-cubes SINFONI de l'échantillon de simulations MIRAGE a été analysé avec l'algorithme développé pour les données MASSIV. Cet outil ajuste un ou plusieurs profils Gaussiens sur chaque spectre du cube considéré et récupère l'information de flux, de position et de largeur de chacune des raies d'émission. À partir de ces informations, la routine d'analyse est en mesure de produire des cartes 2D associées à chacune des quantités : carte de flux, carte de vitesses, carte de dispersion de vitesses.

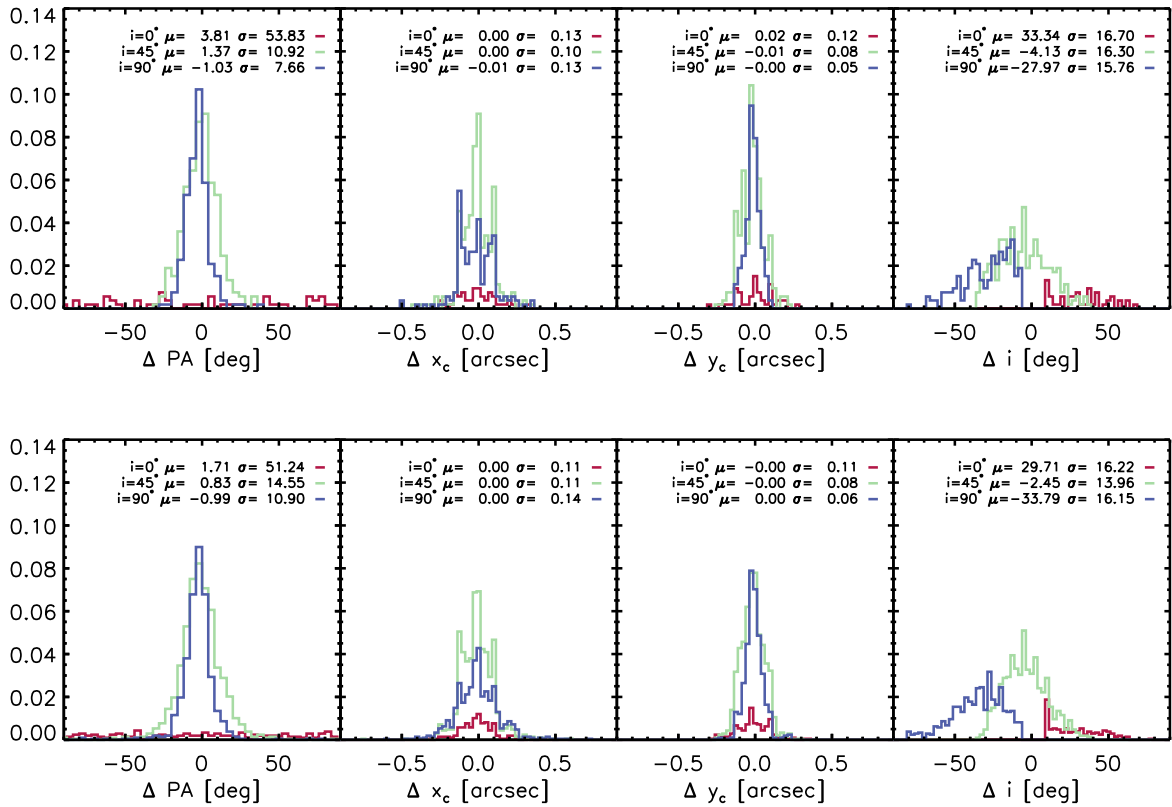


Figure 5.9 – Histogrammes des écarts entre les paramètres géométriques retrouvés par la routine KINEGEN et les paramètres géométriques injectés dans la routine de projection. On mesure respectivement l'écart entre la valeur retrouvée (indexée *rec*) et la valeur injectée (indexée *inj*) $\Delta PA = PA_{ref} - PA_{inj}$, $\Delta x_c = x_{c,rec} - x_{c,inj}$, $\Delta y_c = y_{c,rec} - y_{c,inj}$ et $\Delta i = i_{rec} - i_{inj}$. Pour chaque inclinaison injectée, la médiane μ et la dispersion σ de la distribution sont indiquées. Figure du haut : cas des simulations de disques isolés. Figure du bas : simulation de fusions.

Dans le chapitre 1, j'ai mis en place une analyse cinématique basée sur la méthode de kinétrie. L'analyse de disques exponentiels de la section 1.2.5 m'a permis de définir une procédure permettant d'ajuster les paramètres géométriques (KINEGEN) des galaxies modélisées et d'en extraire par la suite une décomposition en harmoniques sphériques. Je peux donc utiliser maintenant mon échantillon de pseudo-observations issues des simulations MIRAGE comme un échantillon de référence pour quantifier la capacité de cette procédure à

retrouver les paramètres géométriques et physiques de galaxies observées à grand décalage spectral à partir duquel une classification cinématique pourra être mise en place.

Comme la géométrie d'entrée de chacune des pseudo-observations est définie précisément, il est possible d'évaluer la qualité de l'ajustement des paramètres géométriques par la routine `KINEGEN` en utilisant la méthode définie dans la section 1.2.5. La distribution de la différence entre paramètres géométriques d'entrée et paramètres géométriques retrouvés est présentée dans la figure 5.9 (graphe du haut) pour les simulations isolées et dans la figure 5.9 (graphe du bas) pour le cas des simulations de fusions. La statistique élevée des pseudo-observations permet de quantifier

- Le PA est retrouvé avec une précision à 1σ de l'ordre de 10° , ce qui est très satisfaisant vu la basse résolution spatiale des observations par rapport à celle des simulations.
- Le centre est aussi retrouvé avec une précision très satisfaisante, avec une erreur à 1σ de l'ordre de $0.1''$, soit moins de 1 kilo-parsec pour un décalage spectral $z=1.7$.
- L'inclinaison est le paramètre le plus compliqué à ajuster pour les algorithmes de modélisation des champs de vitesses. Le cas $i = 0^\circ$ n'est pas à prendre en compte puisque l'on n'observe aucun mouvement de rotation du disque, ce paramètre est donc nécessairement dégénéré comme le montre l'histogramme des figures 5.9. Cependant, dans le cas $i = 45^\circ$, le pic de la distribution se trouve à 40° avec une dispersion de l'ordre de 15° ce qui donne une précision d'étonnamment bonne qualité et qui confirme la capacité à retrouver les paramètres de rotation déprojetés du plan du disque. Le cas $i = 90^\circ$ est un cas extrême où l'inclinaison retrouvée est plus dispersée, mais statistiquement piquée vers les hautes inclinaisons ce qui est très encourageant.

5.3.2 Estimateurs de perturbation des simulations

Afin de pouvoir déterminer quels paramètres sont pertinents pour estimer le degré de perturbation de chaque pseudo-observation, il est nécessaire de disposer d'estimateurs de perturbation cinématique à la résolution de la simulation numérique. Une corrélation entre cet estimateur "pleine résolution" et un coefficient calculé à partir des harmoniques sphériques obtenue par l'analyse des champs de vitesses `SINFONI` permettrait de pouvoir faire le lien entre des signatures cinématiques à la résolution de la simulation et des signatures cinématiques projetées sur le plan du ciel à la résolution de l'instrument `SINFONI`.

Dans chaque simulation `MIRAGE` contenant n_{cell} cellules AMR, j'ai tout d'abord construit les estimateurs adimensionnels permettant de quantifier le degré d'écart à une rotation purement différentielle du disque. Pour ce faire, je me place dans le référentiel cylindrique centré sur le bulbe stellaire et où le moment angulaire du disque est colinéaire à l'axe z . Pour chaque cellule de gaz i dans le rayon optique du disque stellaire ($r \leq 3.2r_{stars}$), je calcule les trois composante cylindriques $v_{r,i}$, $v_{\phi,i}$, $v_{z,i}$. J'en déduis plusieurs estimateurs construits sur de moyennes pondérées de rapports de vitesses.

L'estimateur

$$E_1 = \frac{\sum_{i=0}^{n_{cell}} m_i \frac{\sqrt{v_{r,i}^2 + v_{z,i}^2}}{v_{\phi,i}}}{\sum_{i=0}^{n_{cell}} m_i} \quad (5.12)$$

quantifie la quantité de mouvement non circulaire normalisée à la vitesse circulaire locale.

L'estimateur

$$E_2 = \frac{\sum_{i=0}^{n_{\text{cell}}} m_i \frac{\sqrt{v_{r,i}^2 + v_{z,i}^2}}{\langle v_\phi \rangle}}{\sum_{i=0}^{n_{\text{cell}}} m_i} \quad (5.13)$$

quantifie la quantité de mouvement non circulaire normalisée à la vitesse circulaire moyenne pondérée par la masse dans le rayon optique du disque.

L'estimateur

$$E_3 = \frac{\sum_{i=0}^{n_{\text{cell}}} m_i \frac{v_{r,i}}{v_{\phi,i}}}{\sum_{i=0}^{n_{\text{cell}}} m_i} \quad (5.14)$$

quantifie la quantité de mouvement radial normalisée à la vitesse circulaire locale.

L'estimateur

$$E_4 = \frac{\sum_{i=0}^{n_{\text{cell}}} m_i \frac{v_{z,i} z_i / |z_i|}{v_{\phi,i}}}{\sum_{i=0}^{n_{\text{cell}}} m_i} \quad (5.15)$$

quantifie la quantité de mouvement vertical d'éloignement au disque (d'où le facteur $z_i/|z_i|$) normalisée à la vitesse circulaire locale.

L'estimateur

$$E_5 = \frac{\sum_{i=0}^{n_{\text{cell}}} m_i \frac{v_{r,i}}{\langle v_\phi \rangle}}{\sum_{i=0}^{n_{\text{cell}}} m_i} \quad (5.16)$$

quantifie la quantité de mouvement radial normalisée à la vitesse circulaire moyenne pondérée par la masse dans le rayon optique du disque.

L'estimateur

$$E_6 = \frac{\sum_{i=0}^{n_{\text{cell}}} m_i \frac{v_{z,i} z_i / |z_i|}{\langle v_\phi \rangle}}{\sum_{i=0}^{n_{\text{cell}}} m_i} \quad (5.17)$$

quantifie la quantité de mouvement vertical d'éloignement au disque normalisée à la vitesse circulaire moyenne pondérée par la masse dans le rayon optique du disque. Dans la figure 5.10, je trace l'évolution de ces six coefficients en fonction du temps de coalescence t_c .

De manière similaire à l'analyse présentée dans Powell et al. (2013), j'ai aussi calculé un coefficient quantifiant la dispersion de vitesse du gaz. Pour ce faire, je considère une grille Cartésienne 3D composée de n_{grid} éléments régulièrement mesurant 100 pc. Dans chaque sous-volume de la grille de dispersion, j'estime le coefficient de dispersion 1D σ_{1D} en sommant quadratiquement la dispersion des composantes xyz :

$$\sigma_{1D} = \sum_{i=0}^{n_{\text{grid}}} \frac{\sqrt{\sigma_{x,1D,i}^2 + \sigma_{y,1D,i}^2 + \sigma_{z,1D,i}^2}}{\sqrt{3}} \quad (5.18)$$

avec $\sigma_{x,1D}$ la dispersion de vitesse pondérée par la masse de la composante Cartésienne x des cellules AMR contenues dans l'élément i de la grille de dispersion. L'évolution de la dispersion de vitesse 1D en fonction du temps de coalescence t_c est présentée dans la figure 5.11.

La mesure ces estimateurs dans chacune des simulations après la coalescence dans le cas des fusions, permet de quantifier l'effet de la fusion sur le support dynamique en fonction du temps. On peut tirer plusieurs conclusions de la figure montrant l'évolution des ces coefficients :

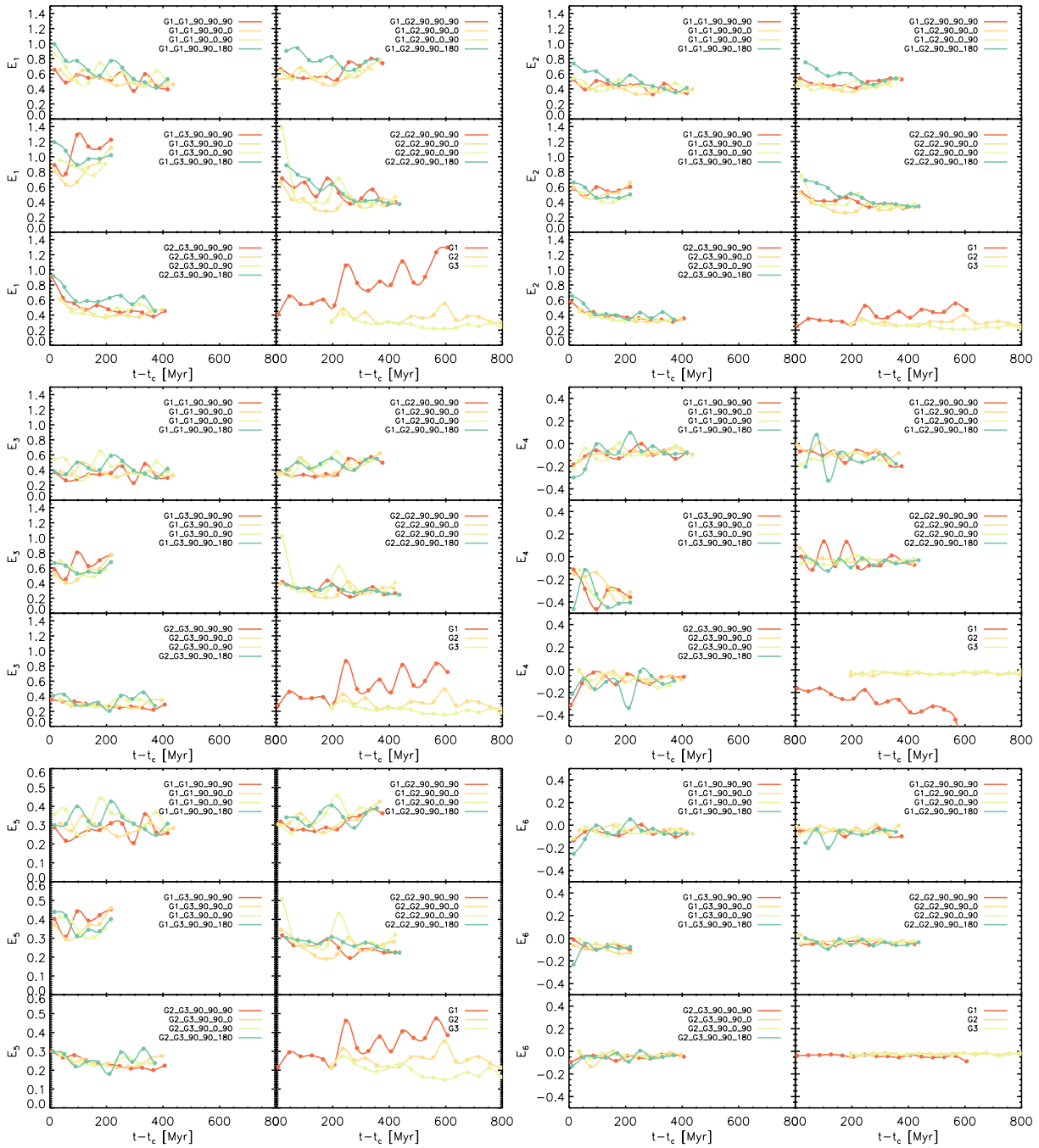


Figure 5.10 – Évolution des estimateurs de perturbation cinématique du gaz E_1, E_2, E_3, E_4, E_5 et E_6 en fonction du temps de coalescence, i.e le temps écoulé depuis la coalescence des progéniteurs. Chaque panneau de chaque figure explore une orientation initiale des disques pour une masse totale donnée. La configuration de chaque simulation MIRAGE est indiquée par la concaténation $G_i-G_j-\theta_1-\theta_2-\kappa$.

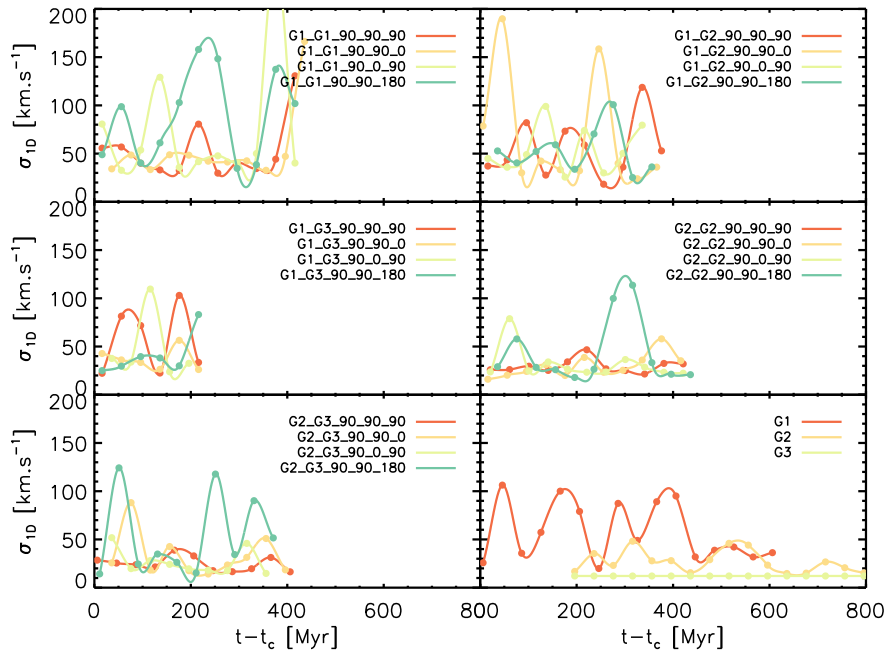


Figure 5.11 – Évolution de l'estimateur de dispersion de vitesse du gaz σ_{1D} en fonction du temps de coalescence, i.e le temps écoulé depuis la coalescence des progéniteurs. Chaque panneau explore une orientation initiale des disques pour une masse totale donnée. La configuration de chaque simulation MIRAGE est indiquée par la concaténation $G_i-G_j-\theta_1-\theta_2-\kappa$.

- La fusion dans les simulations MIRAGE imprime une signature dynamique très modeste comparée au cas des disques isolés. Notamment, la simulation du disque isolé G1 montre une forte instabilité dynamique du fait de la présence de granules très massifs peu affectés par la rétro-action stellaire. Les fusions de ces granules redistribuent une grande quantité de moment angulaire dans le plan du disque et forment un disque tout aussi perturbé que dans les cas de fusion.
- La présence de granules très riches en gaz semble dominer l'évolution à cours terme de la dynamique. Bien que la fusion semble laisser une empreinte dans toutes les simulations à $t - t_c = 0$, le temps de relaxation post-coalescence est très court.
- L'estimateur de dispersion de vitesses du gaz montre une évolution très chaotique, dominée par les éjections de gaz dans le plan du disque dues aux multiples fusions de granules. Seule la simulation du disque isolé G3 montre une dispersion constante au cours du temps ($\sim 20 \text{ km.s}^{-1}$), du fait d'une rétro-action stellaire plus efficace pour détruire des granules bien moins massifs que dans les simulations G2 et G1.

D'après ces estimateurs, il me semble donc difficile dans les simulations MIRAGE de réussir à distinguer un disque en évolution séculaire d'un disque reformé suite à une fusion récente. On peut néanmoins espérer détecter des cas de fusion pré-coalescence, En évaluant la corrélation entre ces estimateurs "pleine résolution"

5.3.3 Perturbation cinématique des pseudo-observations

Afin de quantifier précisément le support dynamique des pseudo-observations MIRAGE, j'ai évalué les coefficients cinématiques définis dans le chapitre 1 en réalisant une décomposition en harmoniques sphériques dont les paramètres géométriques ont été retrouvés par la routine KINEGEN. J'ai souhaité en particulier évaluer la capacité des coefficients de perturbation cinématique v_{asym} et σ_{asym} à retrouver une information de perturbation dynamique dans les simulations à pleine résolution. Dans la figure 5.12, j'évalue en fonction de l'inclinaison d'entrée et du type de simulation (fusion ou disque isolé) les paramètres v_{asym} et σ_{asym} mesurés sur les 4048 pseudo-observations. La couleur de chaque pixel permet d'évaluer le temps moyen écoulé depuis la coalescence dans la région du diagramme délimitée par ledit pixel. On constate dans les cas de fusions que les pseudo-observations situées avant et juste après la coalescence se trouvent dans la partie supérieure droite du diagramme. Dans la figure 5.13, je réalise une figure identique, mais seulement pour les pseudo-observations qui se situent après la coalescence des disques. En comparant à la figure précédente, on peut en déduire que ce sont essentiellement les champs de vitesses pré-coalescence qui viennent peupler cette région du diagramme pour les cas des lignes de visée avec une inclinaison d'entrée du disque de 45 degrés. La majorité des disques observés de face se trouvent dans la région supérieure du diagramme de Shapiro, ce qui est un résultat attendu. On aurait pu néanmoins imaginer détecter plus de rotation dans le cas de disques issus de fusions et vus de face, cas où les parties centrales et externes du disque de gaz peuvent être dynamiquement découplées. Dans la figure 5.15, je montre le même diagramme, mais en affichant cette fois dans chaque pixel la valeur moyenne du coefficient E_2 . On observe pour les pseudo-observations de fusion, inclinées de 45 degrés, une distribution non-aléatoire de cet estimateur de perturbation dynamique, confirmant la capacité des coefficients v_{asym} et σ_{asym} à mesurer une information réelle de perturbation, malgré la PSF instrumentale et le rapport signal à bruit comparable à celui des observations MASSIV.

Dans la figure 5.14, je trace la distribution statistique de la totalité des pseudo-observations, qu'il s'agisse d'une fusion ou d'un disque isolé. La couleur des pixels encode le nombre de pseudo-observations qui se trouvent à l'intérieur de la région que le pixel délimite sur le diagramme. Environ 89% de l'échantillon est détecté dans la région "disque isolé" ($K_{asym} < 0.5$), alors que 87% des simulations modélisent un fusion très riche en gaz. Plus généralement, en utilisant le critère de détection $K_{asym} = 0.5$ de Shapiro et al. (2008) pour classer les fusions et les disques isolés, seulement 352 pseudo-observations sur un total de 3520 pseudo-observations de fusions sont mesurées avec $K_{asym} > 0.5$. Ces 352 pseudo-observations se situent en moyenne au temps de coalescence ($t - t_c \approx 0$ Myr) avec une dispersion de 170 Myr. Il est donc probablement nécessaire de trouver d'autres coefficients de perturbation cinématique qui permettraient de mieux distinguer les cas de fusions des cas de disques isolés et/ou relaxés.

Le travail d'analyse cinématique des pseudo-observations a représenté la dernière étape de ma thèse. L'analyse présentée dans cette section reste encore une ébauche, et a vocation seulement à montrer le potentiel scientifique de l'échantillon de pseudo-observations que j'ai défini. On peut néanmoins tirer quelques conclusions du début de l'analyse des résultats de l'expansion en harmoniques sphériques des champs de vitesses du jeu de pseudo-observations présenté dans ce chapitre.

- Les coefficients de perturbation cinématique v_{asym} et σ_{asym} semblent corrélés au temps $t - t_c$ (temps depuis la coalescence). Cette corrélation se retrouve dans plusieurs esti-

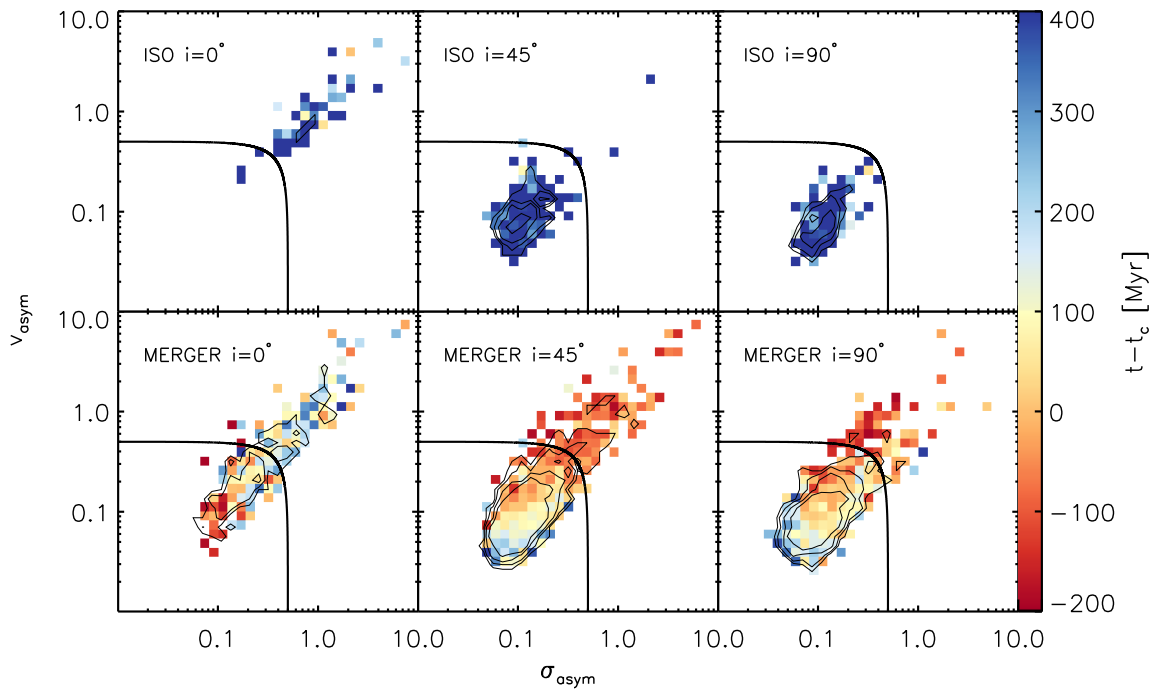


Figure 5.12 – Diagrammes de mesure de l’asymétrie des champs de vitesses et de dispersion de vitesses comme défini par Shapiro et al. (2008). Le rang du haut présente les diagrammes d’asymétries pour les simulations de disques isolés, tandis que le rang du bas présente les diagrammes d’asymétries pour les simulations de fusions. Sur chacun des six diagrammes est indiquée l’inclinaison injectée dans les pseudo-observations. La couleur de chaque pixel d’un diagramme permet de visualiser l’âge de coalescence moyen dans la région du diagramme qu’il occupe. Les contours permettent de visualiser la distribution statistique des pseudo-observations sur les différents diagrammes d’asymétries et sont tracés pour les niveaux [2,4,8,16] pseudo-observations.

mateurs de perturbation dynamique mesuré à pleine résolution dans le rayon optique du disque observé, notamment E_2 montré dans la figure 5.15.

- La frontière $K_{asym}=0.5$ ne permet pas de distinguer les disques isolés des cas de fusion dans les pseudo-observations de l’échantillon MIRAGE, dont la grande majorité réside dans une région très restreinte du diagramme de Shapiro.
- Les seules pseudo-observations qui se positionnent au dessus de la limite $K_{asym} = 0.5$ sont (i) les disques vu de face, (ii) les disques qui sont le résultat immédiat de la coalescence et (iii) les disques en train de finir leur fusion et dont le champ de vitesses ne permet pas de distinguer deux composantes distinctement. Les galaxies qui se trouvent au-delà de la limite $K_{asym} = 0.5$ sont pour la plupart des objets dont on observe seulement les mouvements verticaux du gaz.

Le travail présenté dans ce chapitre permet d’envisager plusieurs pistes afin de permettre

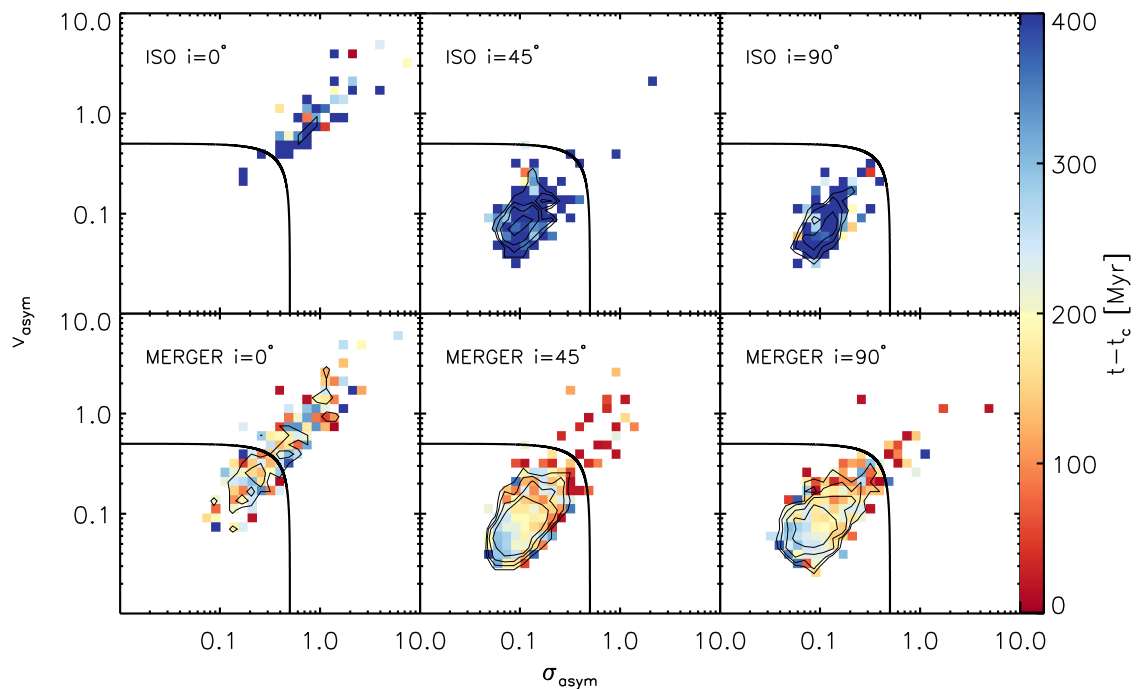


Figure 5.13 – Diagrammes de mesure de l’asymétrie des champs de vitesses et de dispersion de vitesses comme défini par Shapiro et al. (2008), pour les pseudo-observations post-coalescence seulement. Le rang du haut présente les diagrammes d’asymétries pour les simulations de disques isolés, tandis que le rang du bas présente les diagrammes d’asymétries pour les simulations de fusions. Sur chacun des six diagrammes est indiquée l’inclinaison injectée dans les pseudo-observations. La couleur de chaque pixel d’un diagramme permet de visualiser l’âge de coalescence moyen dans la région du diagramme qu’il occupe. Les contours permettent de visualiser la distribution statistique des pseudo-observations sur les différents diagrammes d’asymétries et sont tracés pour les niveaux [2,4,8,16] pseudo-observations.

de quantifier la perturbation cinématique des galaxies dans les observations spectroscopiques à grand décalage spectral. Le jeu de pseudo-observations permet d’envisager la création de nouveaux coefficients basés sur les harmoniques sphériques. Bien que la détection de signatures spécifiques aux fusions semble une tâche difficile, il est envisageable de comparer ce travail à des pseudo-observations basées sur des simulations moins riches en gaz, afin de déterminer si les fusions impriment des signatures plus fortes que l’évolution séculaire à plus faible fraction de gaz du disque. Finalement, une piste à explorer serait de comparer observations et simulations sans perdre l’information spatiale : comparer les ellipses de la méthode de kinémetrie unes à unes, et regarder dans quelles pseudo-observations on détecte des coefficients ayant un distribution radiale similaire pourrait éventuellement permettre d’affiner la technique de détection de fusion par l’analyse cinématique.

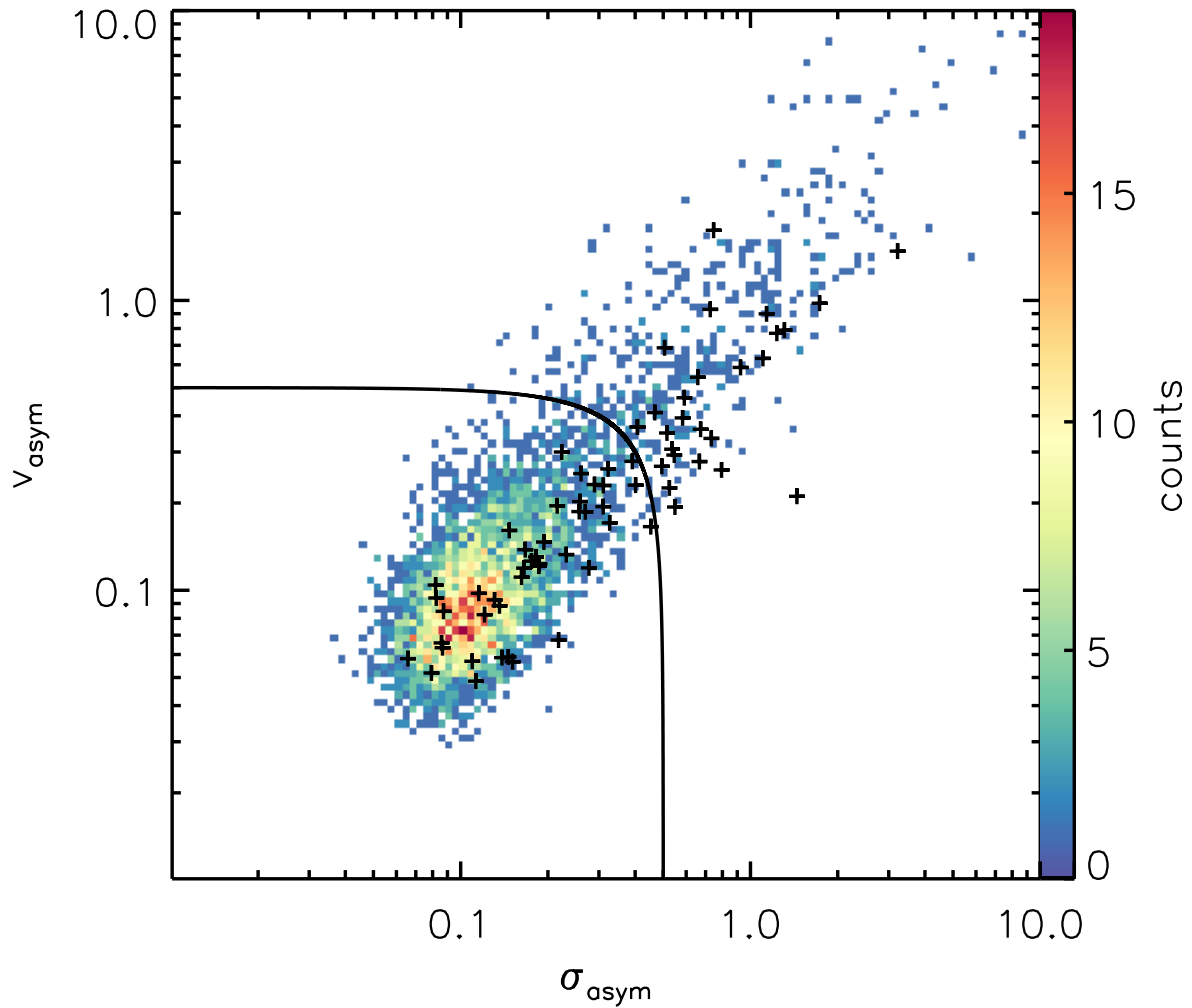


Figure 5.14 – Diagramme de mesure de l'asymétrie des champs de vitesses et de dispersion de vitesses comme définie par Shapiro et al. (2008), pour l'échantillon MASSIV (symbole "+"). La distribution statistique de l'échantillon MIRAGE est superposée, grâce à une échelle de couleur.

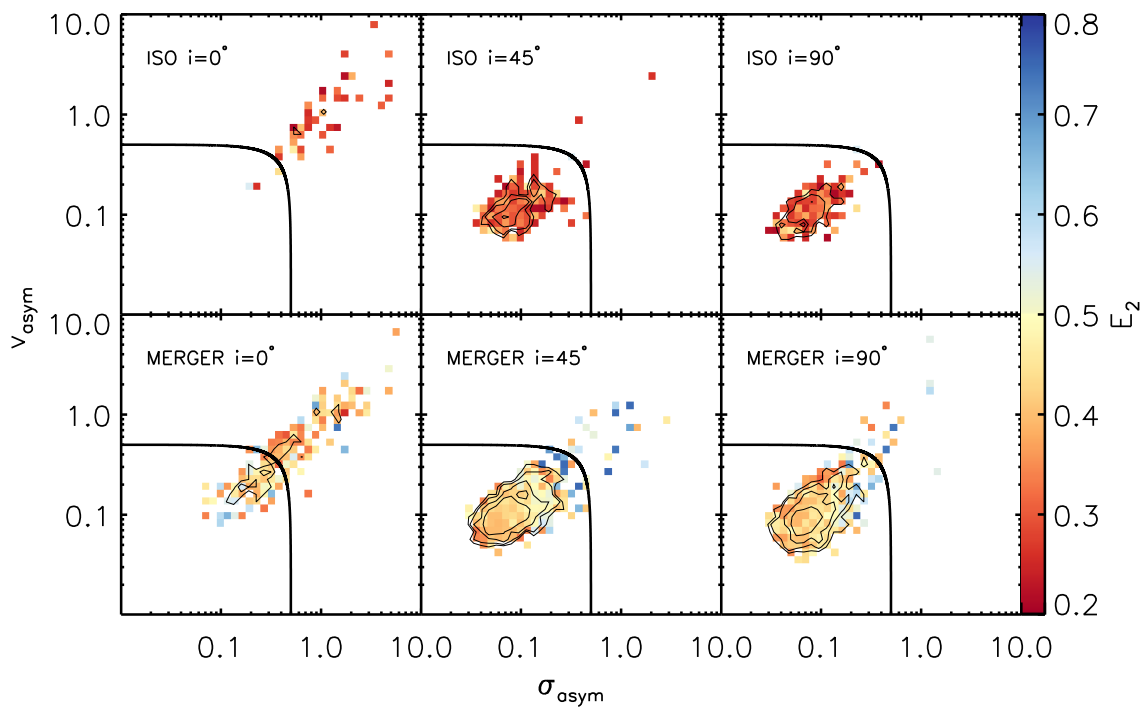


Figure 5.15 – Diagrammes de mesure de l’asymétrie des champs de vitesses et de dispersion de vitesses comme définie par Shapiro et al. (2008). Le rang du haut présente les diagrammes d’asymétries pour les simulations de disques isolés, tandis que le rang du bas présente les diagrammes d’asymétries pour les simulations de fusions. Sur chacun des six diagrammes est indiquée l’inclinaison injectée dans les pseudo-observations. La couleur de chaque pixel d’un diagramme permet de visualiser l’âge de coalescence moyen dans la région du diagramme qu’il occupe. Les contours permettent de visualiser la distribution statistique des pseudo-observations sur les différents diagrammes d’asymétries et sont tracés pour les niveaux [2,4,8,16] pseudo-observations.

Conclusion & Perspectives

Les observations spectroscopiques des galaxies de l'Univers lointain apportent une quantité considérable d'informations qui permet de mieux contraindre les scénarios d'évolution des galaxies. Cependant, de nombreuses questions persistent : Quand les galaxies se sont-elles formées ? Quand les premières étoiles ont-elles été formées ? Quand apparaissent les premiers disques stellaires ? Comment s'est construite la séquence de Hubble moderne ? Dans quelles mesures les fusions successives de galaxies et l'accrétion de gaz primordial sur les disques participent-elles au grand assemblage de masse cosmologique ? L'analyse morphologique des galaxies lointaines ($z > 1$) révèle que la majorité des galaxies spirales a dû passer par une phase d'instabilité gravitationnelle où les disques stellaires et gazeux étaient dominés par plusieurs granules concentrant une grande fraction de la masse baryonique. Cette phase granulaire est actuellement partiellement modélisée par les simulations cosmologiques. Celles-ci ne possèdent pas encore les moyens d'atteindre la résolution numérique permettant d'échantillonner correctement la structure verticale des disques, et donc de simuler leur fragmentation de manière réaliste. Cette thèse a pour objectif d'apporter de nouveaux éléments quant à l'évolution de ces disques "juvéniles", notamment lorsqu'ils subissent une fusion. Mon travail s'est articulé sur deux axes complémentaires : le premier est observationnel et se concentre sur l'analyse cinématique des galaxies jeunes, le second est numérique et vise à simuler les galaxies jeunes très riches en gaz et hautement instables.

Observations

Cette thèse a été l'occasion de m'impliquer dans la réduction et l'analyse des données de l'échantillon spectroscopique MASSIV, qui regroupe 83 galaxies entre $1 < z < 2$, époque où l'on observe un pic dans le taux de formation stellaire cosmologique. Après avoir réduit quelques données, ma première contribution au projet MASSIV a consisté à mettre au point une méthode d'analyse des champs de vitesses permettant de quantifier et qualifier le support rotationnel de chacune des galaxies. J'ai développé un programme, dénommé KINEGEN combinant une décomposition du champ de vitesses en harmoniques sphériques⁵ à l'aide d'un algorithme génétique qui se révèle très efficace pour éviter les nombreux minimum locaux dans la fonction de χ^2 . J'ai mis le programme KINEGEN à disposition de la communauté scientifique librement⁶. L'ajustement de l'inclinaison d'un disque à partir de données cinématiques est de loin le processus le plus complexe, essentiellement à cause de la faible résolution spatiale des observations de galaxies à grand décalage spectral. Afin de tenir compte des effets de lissage induits par la réponse impulsionnelle spatiale du télescope, j'ai construit environ 200 modèles de champs de vitesses de disques ayant un profil de masse exponentiel et explorant les paramètres d'inclinaison, de taille du disque et de réso-

5. aussi désigné par le terme *kinemetry* en anglais

6. <http://code.google.com/p/kinegen/>

lution spatiale. À partir de ces modèles, j'ai trouvé une expression analytique qui détermine la forme optimale des ellipses le long desquelles sont extraits les coefficients harmoniques, ce qui permet de retrouver de manière plus précise le paramètre d'inclinaison du disque. J'ai ajusté les différents paramètres géométriques, puis j'ai réalisé la décomposition en harmoniques sphériques de l'échantillon MASSIV ayant servi à l'extraction des paramètres cinématiques des galaxies. À partir des informations cinématiques obtenues, j'ai réalisé une classification cinématique semblable à celle réalisée pour l'échantillon SINS. Je trouve une fraction de 30% de galaxies susceptibles d'être une fusion en cours. Enfin, j'explore différentes pistes concernant la construction de nouveaux coefficients permettant de quantifier le degré de perturbation d'un champ de vitesses.

Simulations

Un moyen de contraindre les scénarios d'évolution des galaxies est de recourir à des simulations numériques. Dans le cadre de cette thèse, je me suis cantonné au cas des simulations dites "idéalisées" pour lesquelles les galaxies évoluent au sein d'un environnement fermé, en dehors de tout contexte cosmologique. L'avantage des simulations idéalisées est de permettre d'atteindre des résolutions physiques bien plus élevées que celles de simulations cosmologiques, qui elles, au contraire, modélisent un sous-volume d'Univers dont les paramètres sont fixés grâce à la cosmologie observationnelle. J'ai écrit un logiciel permettant la construction de conditions initiales proches de l'équilibre dynamique. Le logiciel nommé DICE est le produit de ce travail, je l'ai rendu disponible librement à la communauté scientifique (code.google.com/dice-project). DICE est capable de générer une grande variété de modèles de galaxies à l'équilibre du fait de sa large paramétrabilité. DICE est aussi en mesure de placer deux galaxies sur une trajectoire képlérienne afin de simuler une rencontre galactique. L'effort apporté à l'écriture du code et à la facilité d'accès et d'utilisation par des utilisateurs tiers en feront, je l'espère, un outil largement utilisé dans les prochaines années.

Disposer d'un outil de construction de conditions initiales de disque galactique m'a permis de construire l'échantillon MIRAGE, composé de 20 simulations de fusions et 3 simulations de disques isolés dans les conditions de l'Univers à $z=2$. Les conditions initiales des simulations de fusions explorent quatre orientations initiales distinctes des disques pour cinq configurations différentes de masse. Tous les disques possèdent des fractions de gaz élevées (~60%) et sont continuellement alimentés par une accrétion idéalisée de gaz chaud du milieu inter-galactique qui est réalisée en maintenant une densité minimale dans la grille de simulation.

Les résultats de l'analyse de l'échantillon MIRAGE sont listés ci-dessous :

1. Les granules⁷ formés par instabilités gravitationnelles dans le disque primordial sont en mesure de survivre plusieurs centaines de millions d'années, durant lesquelles ils migrent vers le centre de masse du système grâce à un mécanisme de friction dynamique gravitationnelle. Malgré une importante éjection continue de gaz, ces granules sont capables de maintenir une masse constante du fait qu'ils se déplacent dans un milieu interstellaire très turbulent duquel ils accrètent suffisamment de masse pour compenser les éjections.

7. "clumps" en anglais.

2. L'implémentation des mécanismes de rétro-action stellaire des étoiles de type OB par pression radiative et par injection d'énergie thermique des supernovae de type II génère des flots⁸ massifs de gaz sortant des granules formés par instabilités de Jeans. Ce taux d'éjection est aux alentours du taux de formation stellaire dans les granules. Ces flots possèdent des propriétés en accord avec les observations.
3. L'accrétion continue de gaz par les granules permet de maintenir le taux de formation stellaire à un niveau relativement constant durant plusieurs centaines de millions d'années. L'âge moyen des étoiles présentes dans un granule est significativement plus faible que l'âge même du granule. Les étoiles qui sont formées dans le granule sont expulsées graduellement sous l'effet des forces de marées qui sont à l'oeuvre lors d'interactions entre granules, de telle sorte que la moitié des étoiles formées *in situ* quitte le granule 200 millions d'années plus tard. Ces effets des forces de marée permettent de garder une population stellaire relativement jeune dans un granule dont l'âge de formation peut remonter de plusieurs centaines de millions d'années.
4. Les taux de formation stellaire globaux des disques des simulations de l'échantillon MIRAGE fluctuent fortement, avec une dispersion proche de 30%. La formation stellaire s'effectue principalement dans les complexes granulaires et se trouve fortement stimulée lorsque ces granules se rencontrent et fusionnent.
5. Les simulations de fusion de galaxies de l'échantillon MIRAGE ne sont pas en mesure d'augmenter l'efficacité de formation stellaire de manière significative au-delà de la dispersion intrinsèque du taux de formation stellaire. Ce résultat surprenant va à l'encontre de l'idée répandue selon laquelle la fusion de deux disques gazeux induit une flambée de formation stellaire. Bien que divers mécanismes puissent être à l'origine de la disparition de la flambée de formation stellaire dans les fusions de galaxies juvéniles, une explication possible serait que le haut niveau de fragmentation et de turbulence du gaz pourrait être en mesure de saturer la formation stellaire. Dans les simulations MIRAGE, la fragmentation et la turbulence du gaz sont maintenues à un niveau élevé par l'injection d'énergie des étoiles jeunes dans le milieu interstellaire ainsi que par l'interaction de granules dont la haute fraction de gaz est maintenue par une accrétion continue provenant du halo chaud circum-galactique. Ce résultat, qui est à mettre en parallèle avec la remarquable homogénéité du taux de formation stellaire spécifique des galaxies à grand décalage spectral (Elbaz et al., 2007, 2011 ; Nordon et al., 2012), suggère que l'accroissement de l'efficacité de formation stellaire est rare dans les fusions de galaxies jeunes. Une flambée de formation stellaire induite par une fusion pourrait donc être un mécanisme qui requerrait des conditions spécifiques peu fréquentes dans l'Univers jeune pour provoquer une telle augmentation d'efficacité de formation stellaire.
6. Les simulations de l'échantillon MIRAGE reproduisent une évolution de la taille du disque stellaire en accord avec les relations masse-taille calibrées sur les galaxies locales. L'importante redistribution du moment angulaire, notamment par les granules riches en gaz, permet aux disques d'accroître leur taille continuellement, et ce, malgré la présence d'une fusion. Les fusions de galaxies sont pourtant un processus considéré comme efficace pour concentrer la matière baryonique et non-baryonique vers le centre de masse. Les simulations étudiées dans cette thèse démontrent que l'accrétion cosmologique de gaz froid transportant un fort moment angulaire n'est donc pas une condition *sine qua non* pour construire radialement les disques des galaxies spirales.

8. "outflows" en anglais

7. L'accrétion de gaz chaud virialisé telle qu'elle est implémentée dans les simulations MIRAGE est très efficace pour alimenter le disque en gaz et compenser à tout moment la consommation par la formation stellaire, de telle manière que la fraction de gaz se maintient tout au long de l'évolution des simulations. Les granules sont capables d'accréter le gaz environnant et de maintenir une fraction de gaz supérieure à 90% pour les plus massifs d'entre eux. L'interaction de granules extrêmement riches en gaz est capable d'éjecter de fortes quantités de gaz par transfert de moment angulaire vers des queues de marée intra-disque.
8. La fusion des complexes granulaires riches en gaz est donc capable de conduire plusieurs processus d'évolution de la galaxie hôte qui permettent de voir sous un regard nouveau différentes problématiques actuelles concernant la distribution de masse dans les simulations numériques :
 - Éjection de gaz - les queues de marées formées par la coalescence de deux ou plusieurs granules sont capables d'emporter de grandes quantités de masse en dehors de la galaxie hôte. Les taux d'éjection sont comparables aux taux de formation stellaire dans les cas les plus massifs. Les éjections, qui prennent la forme de bras spiraux avec des vitesses radiales élevées ($>500 \text{ km.s}^{-1}$), se situent toujours dans un plan quasiment co-planaire au disque, puisque qu'elles résultent du transfert de moment angulaire de deux granules dont le propre moment angulaire est quasi-co-linéaire à celui du disque. Ce mécanisme, bien qu'atypique, pourrait être une réponse aux problèmes de sur-concentration des baryons produits dans les disques des simulations numériques cosmologiques. Les simulations qui arrivent à enlever suffisamment de baryons grâce à la rétro-action stellaire forment systématiquement des galaxies elliptiques et n'arrivent pas à former suffisamment de galaxies spirales. En revanche, le mécanisme présenté ici est capable d'enlever une grande quantité de baryons du disque, sans pour autant en altérer l'existence. Une fraction des matériaux éjectés retombera ultérieurement dans le plan du disque, de manière similaire à l'accrétion de gaz froid primordial.
 - Éjection d'étoiles - Les queues de marée intra-disque sont capables d'emporter avec elles une partie des étoiles initialement présentes dans les complexes granulaires. La coalescence des granules est donc un mécanisme efficace pour retirer les étoiles du bulbe stellaire grâce aux forces de marées. De manière équivalente, les granules très riches en gaz qui sont capables d'emmener de grandes quantités de matière vers le centre de la galaxie sont aussi capables d'évacuer une grande partie de ces matériaux vers des régions plus externes. Les étoiles éjectées participent alors à la construction d'un sphéroïde stellaire, composé d'étoiles ayant des orbites significativement plus excentriques que celles du disque. Ce mécanisme permet de ralentir l'accroissement de masse du bulbe stellaire, qui est un problème récurrent dans les simulations cosmologiques. La sur-concentration de masse dans le bulbe peut produire des courbes de rotation très piquées dans les régions centrales des galaxies qui ne sont pas observées dans l'Univers local.
 - Matière sombre - Les simulations de l'échantillon MIRAGE sont toutes initialisées avec un profil de matière sombre piqué au centre. Ce profil, typique dans les simulations cosmologiques, n'est pourtant pas celui déduit par la modélisation de masse des galaxies de l'Univers local, dont les halos semblent présenter un profil de masse plat dans les parties centrales. De nombreux travaux récents étudient

les différentes pistes qui pourraient expliquer les mécanismes à l'origine d'une transformation du profil du halo depuis un profil de masse piqué au centre vers un profil plat. Les mécanismes de rétroaction stellaires sont capables de transformer la structure du halo pour produire un noyau⁹, mais cela n'est valable que pour les galaxies les moins massives (galaxies naines). L'interaction des granules dans les simulations de l'échantillon MIRAGE est capable de chauffer violemment le noyau du halo. Ce processus, qui peut se produire plus d'une dizaine de fois en 1 milliard d'années dans les simulations de disques les plus massifs, est donc en mesure de redistribuer la matière sombre qui se trouve à l'intérieur du kiloparsec central. Ce cas de figure nécessite que les granules formés par instabilités gravitationnelles puissent survivre à l'injection d'énergie par les étoiles jeunes.

Pseudo-observations

Ce projet de thèse s'est initialement construit sur l'idée de comparer des simulations numériques à des données d'observations spectroscopiques afin de déterminer s'il était possible de détecter des signatures cinématiques de fusion dans les observations de galaxies lointaines, où la résolution spatiale limite fortement les capacités d'analyse. Pour établir cette comparaison, j'ai développé au cours de cette thèse un programme nommé *Aurora* permettant de reproduire des observations à partir de simulations numériques de galaxies pour un décalage spectral donné. Ce programme est lui aussi librement accessible¹⁰. À partir des simulations MIRAGE, j'ai créé un jeu de plus de 4000 pseudo-observations¹¹ SINFONI de galaxies projetées au décalage spectral $z=1.7$ en simulant des caractéristiques instrumentales de l'instrument SINFONI (champ du vue, échantillonnage spatial et spectral, etc.) et des conditions d'observation (turbulence atmosphérique), en explorant différentes géométries de projection et différents pas de temps. J'ai par la suite analysé les pseudo-observations avec les logiciels de réduction développés pour les données MASSIV de manière à disposer de données strictement comparables aux observations. J'ai ainsi produit des cartes de flux stellaires et $H\alpha$, des champs de vitesses et de dispersion de vitesses comparables aux observations. Les paramètres géométriques de chaque champ de vitesses issus des pseudo-observations ont été ensuite déterminés en utilisant le programme *KINEGEN*. En réalisant une décomposition en harmoniques sphériques grâce à la méthode *kinemetry*, je dispose d'une statistique conséquente qui peut me permettre de déterminer s'il existe des signatures spécifiques de fusions de disques gazeux dans les simulations MIRAGE dans les coefficients de l'expansion en harmoniques sphériques.

Je résume ci-dessous les résultats de cette analyse :

1. Validation de la méthode d'ajustement des paramètres géométriques : la méthode *KINEGEN* est statistiquement capable de retrouver les bons paramètres géométriques des simulations projetées (angle de position de l'axe majeur de la galaxie, coordonnées du centre, inclinaison du disque). Malgré la cinématique complexe du gaz, la méthode développée au cours de cette thèse se révèle être un outil robuste, capable de traiter automatiquement un échantillon de plusieurs milliers de champs de vitesses sans avoir à fixer des paramètres d'ajustement différents pour chaque galaxie.

9. "core" en anglais. Ce terme désigne la région centrale du halo de matière sombre où le profil radial de densité est plat.

10. <http://aurora-py.googlecode.com>

11. "mock observations" en anglais.

2. Détection statistique de fusion : la fusion de galaxie semble être relativement difficile à détecter par l'analyse cinématique des pseudo-observations MIRAGE. Le disque de gaz se reforme très rapidement après la coalescence des deux disques progéniteurs, et la structure fragmentée du gaz laisse peu de signatures cinématiques claires contrairement aux fusions de l'Univers local. Plus généralement, en utilisant le critère de détection $K_{asym} = 0.5$, seulement 352 pseudo-observations sur un total de 3520 pseudo-observations de fusions sont mesurées au dessus de $K_{asym} = 0.5$. Ces 352 pseudo-observations se situent en moyenne au temps de coalescence $(t - t_c) = 0$ Myr avec une dispersion de 170 Myr. Ce résultat confirme la dépendance au modèle numérique utilisé pour "entraîner" une méthode de classification automatique. La détection devient plus claire en phase pre-coalescence, où des effets d'alignements géométriques peuvent cacher une galaxie derrière l'autre. Globalement, la plupart des observations montrent une cinématique peu perturbée duquel il est difficile de détecter la récente présence d'une fusion. L'analyse commencée dans cette thèse demande à être étendue, notamment en cherchant des signatures plus claires dans de nouveaux coefficients construits à partir de la méthode kinemetry. L'analyse d'estimateurs de perturbation dynamique à la résolution de la simulation montre néanmoins qu'il n'est pas envisageable d'obtenir une signature évidente de fusion, du fait du haut niveau de perturbation pré-existant dans les disques isolés. Des simulations complémentaires avec une fraction de gaz plus faible permettraient de voir comment varient ces résultats dans un environnement moins turbulent.

Perspectives

Dans cette thèse, j'ai abordé la problématique de la fusion des galaxies jeunes en faisant le lien entre l'aspect observations spectroscopiques et simulations numériques. La récente mise en service de l'instrument KMOS (Sharples et al., 2013), qui est composé de 24 spectrographes intégral de champ va apporter une quantité considérable de données cinématiques à grand décalage spectral qu'il faudra analyser. La méthode d'analyse cinématique présentée dans cette thèse est à même de traiter de grandes quantités de données parce qu'elle est automatisée et robuste. La base de données de pseudo-observations bâtie au cours de cette thèse pourra en premier lieu être comparée à d'autres échantillons spectroscopiques, et particulièrement aux futures données des instruments KMOS du VLT, MOSFIRE sur le télescope Keck (McLean et al. 2012) et autre radio télescope millimétrique ALMA. Ce dernier devrait d'ailleurs nous éclairer sur le contenu moléculaire des galaxies nécessaire à la compréhension des mécanismes de formation stellaire.

Réaliser des simulations numériques implique nécessairement de faire des choix quant à l'implémentation de la physique. Des simulations très riches en terme de modélisation de processus physiques impliquent de maîtriser un cas plus simple et moins coûteux en temps de calcul, ce qui peut déjà se révéler un travail conséquent. Les simulations MIRAGE ont soulevé des questions passionnantes, que je compte étudier afin de compléter le travail présenté dans ce manuscrit.

Notamment, l'inefficacité des fusions des simulations de disques riches en gaz pour former des étoiles appelle à la réalisation de simulations complémentaires afin de contraindre de manière claire les principaux mécanismes responsables de cette saturation de la formation stellaire. Le rôle de la coalescence des granules riches en gaz dans l'évolution des propriétés physiques des simulations MIRAGE nécessite aussi d'être étudié plus en détail. On pourrait ainsi chercher à déterminer à partir de quelle fraction de gaz la fusion de granules

devient-elle inefficace pour redistribuer la masse baryonique et non baryonique au centre des galaxies. Il serait en particulier intéressant de déterminer s'il est possible dans les simulations cosmologiques d'obtenir un environnement qui serait propice à l'apparition de granules très riches en gaz dans le plan du disque galactique. Dans les simulations MIRAGE, j'ai volontairement omis d'intégrer le traitement physique des trous-noirs super-massifs et de la rétro-action des AGN par souci de gain de temps de calcul. Les AGN sont en mesure de pousser le halo de gaz proche du disque, et donc d'affecter le taux global d'accrétion de gaz. Une question à laquelle il serait nécessaire de répondre concerne l'effet que les AGN pourraient avoir sur les granules issus d'instabilités gravitationnelles. Ces granules seraient-ils toujours en mesure d'accréter autant de gaz malgré l'énergie thermique que l'AGN transférerait au halo de gaz ? Ce mécanisme ne forcerait-il pas l'accrétion du gaz dans le plan galactique perpendiculaire au cône de l'AGN ? Récemment, le code `Ramses` a intégré un module complet de transfert radiatif (Rosdahl et al., 2013) qui permet de suivre l'ionisation des éléments du milieu interstellaire de manière plus réaliste que l'approximation de la sphère de Strömgen que j'ai utilisé dans les simulations MIRAGE. Réaliser ce type de simulations est une amélioration naturelle à apporter aux simulations de l'échantillon MIRAGE. Plusieurs questions fondamentales se posent : les granules sont-ils capables de résister à la pression radiative des étoiles jeunes modélisées par un transfert radiatif complet ? Où se trouve le gaz ionisé dans ces galaxies jeunes extrêmement turbulentes ? Ce type de simulation permettrait notamment de construire des pseudo-observations plus réalistes : on disposerait alors de la fraction d'ionisation en tout point du volume simulé. On peut se demander si l'inclusion du transfert radiatif serait en mesure de modifier la signature cinématique du disque simulé et du disque perçu au travers d'une pseudo-observation. Enfin, réaliser des pseudo-observations avec un post-traitement d'une simulation en incluant un transfert radiatif permettrait de quantifier notre capacité d'interprétation des résultats du cas optiquement fin traité dans cette thèse. Plus particulièrement, l'étude de la largeur des raies d'émission, qui ne peut pas être entreprise sans un traitement radiatif du milieu, permettrait de mieux comprendre la nature de l'augmentation de la turbulence dans les disques jeunes.

Grâce à l'arrivée, sur les télescopes de la classe des 8-10 mètres, des instruments de seconde génération permettant de réaliser des observations de champ intégral, nous avons peu à peu découvert depuis 7 ans combien les galaxies étaient différentes par le passé. Des sondages ont permis d'étudier des galaxies dont les décalages spectraux s'étendent de $z \sim 0.3$ à 3.5, couvrant $\sim 85\%$ de l'âge de l'Univers, ils ont permis de montrer leur évolution dynamique. Le principal problème observationnel est lié à la résolution spatiale limitée de ces données qui permet difficilement d'étudier les galaxies massives et quasiment pas les galaxies de plus faible masse. L'optique adaptative offre de nouvelles perspectives mais les limitations en termes de flux vont rendre la tâche difficile. L'arrivée dans une petite dizaine d'années des ELT et du JWST et ultérieurement encore du radiotélescope SKA, devrait fournir des observations de galaxies à $z > 1$ avec une résolution spatiale comparable à celle dont nous disposons pour les galaxies de l'Univers local avec les plus grands télescopes actuels et lever encore bien des voiles. D'ailleurs, seuls des télescope de la classe des 25-40 mètres auront la sensibilité qui permettra d'étudier la dynamique des étoiles de ces galaxies.

En attendant, beaucoup de travaux pourront néanmoins être réalisés à l'aide des nouveaux moyens d'observation actuels qui correspondent, en quelque sorte, à une troisième génération d'instruments sur les grands télescopes optiques/infrarouge ainsi que grâce à une nouvelle fenêtre dans le domaine submillimétrique offerte par ALMA et sa montée progressive en sensibilité et en résolution (par l'adjonction de nouvelles antennes) qui va certainement contribuer à révolutionner notre vision de la formation et de l'évolution des galaxies

dans les cinq prochaines années (sans oublier le réseau submillimétrique NOEMA dans l'hémisphère nord). Combiné à notre meilleure compréhension des mécanismes physiques et à l'augmentation régulière des capacités des super-calculateurs, de nombreux travaux passionnants sur l'étude de la formation et de l'évolution des galaxies nous attendent dans les années à venir. À ne pas en douter, les questions principales auxquelles il faudra répondre de façon statistique à partir d'échantillons représentatifs sont parmi les suivantes :

- Quelle est la fonction de masse initiale à grand décalage spectral ? Les lois et mécanismes de formation stellaire locaux établis dans l'Univers local pour des disques fins sont-ils toujours valables lorsque l'échelle de hauteur, la densité et la pression du gaz sont beaucoup plus élevées alors que sa métallicité est plus faible (validité du facteur de conversion CO) ? Existe-t-il des seuils ou des effets de saturation de la densité sur la formation stellaire comme nous l'observons dans les simulations MIRAGE ?
- Comment varie le taux de fusion entre galaxies au cours du temps, aux époques où les galaxies étaient globalement beaucoup plus gazeuses qu'elles ne le sont aujourd'hui ? Quelle est l'influence de la distribution des orbites, de l'orientation relative des disques, de leur rapport de masse sur le taux de fusion ? Comment déduire ces paramètres d'un contexte cosmologique complet ? Comment interpréter les signatures cinématiques souvent complexes que l'on observe dans une fraction importante de galaxies alors que, comme nous l'avons vu dans l'analyse de l'échantillon MIRAGE, la signature d'une fusion entre deux systèmes très gazeux est un phénomène transitoire ? Comme nous l'avons également entrevu dans ce travail, dispose-t-on d'une résolution angulaire suffisante pour interpréter les écarts observés à la rotation circulaire ?
- Comment varie la fraction de disques à chaque époque pour chaque intervalle de masse de galaxies ? L'étude de la dispersion dans les relations de Tully-Fischer et plus généralement des lois d'échelle, est intimement liée à la physique et aux devenir même de ces disques. Ces disques sont-ils les précurseurs des galaxies actuelles (elliptiques, spirales, lenticulaires) ou pas ?
- Pourquoi observe-t-on une fraction aussi importante de galaxies gazeuses et formant beaucoup d'étoiles, supportées gravitationnellement par la dispersion de vitesse du gaz et des étoiles nouvellement formées ? Ces galaxies n'ont pas d'analogue dans l'Univers local. Sont-elles à l'équilibre dynamique ? Sont-elles le résultat de fusions ? Sont-elles alimentées par des flots continus de gaz cosmique entrant ? Existe-t-il des flots importants de gaz sortant du disque de ces galaxies ? Vers quoi vont-elles évoluer ?
- Quelle est la physique des granules ? Ces granules sont-ils vraiment résolus ou bien sont-ils constitués par des entités plus petites liées (ou pas) par un support de pression gravitationnelle (dispersion de vitesses) ? Ont-ils atteint une masse de Jeans qui leur permet de suivre une évolution propre dans le potentiel galactique ? Comment se forment-ils, comment évoluent-ils, quelle est leur durée de vie, comment sont-ils détruits et/ou reformés ? Quel est leur taux de migration vers le centre de la galaxie ? Sont-ils régulièrement alimentés en gaz frais cosmique, galactique, ou bien circumgalactique ? Quelle fraction de matière rejettent-ils dans le milieu interstellaire et intergalactique, par quels mécanismes ? Quel est le rôle des métaux et de la poussière ? Contiennent-ils uniquement de la masse baryonique ? Quelle est leur fraction de gaz

moléculaire ? Et comme ils n'ont pas d'équivalent dans l'Univers actuel, comment disparaissent-ils ?

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UNIVERSITÉ D'AIX-MARSEILLE

ÉCOLE DOCTORALE DE PHYSIQUE ET SCIENCES DE LA MATIÈRE

FACULTÉ DES SCIENCES

LABORATOIRE D'ASTROPHYSIQUE DE MARSEILLE

THESE DE DOCTORAT

Discipline : Sciences de la Terre et de l'Univers, Espace

Spécialité : Astrophysique et Cosmologie

Valentin PERRET

Fusions de galaxies juvéniles
Des simulations aux observations

ANNEXES

Soutenue publiquement le 30 janvier 2014

Introduction

Ces annexes regroupent les différents documents ne pouvant pas rentrer dans le corps du texte du fait de leur taille, et/ou de leur contenu trop technique. J'ai notamment produit de nombreuses cartes des simulations et des simulations numériques qui sont très volumineuses. Dans une première annexe, le lecteur pourra trouver les cartes qui n'ont pas pu être présentée dans le corps de la thèse. Les figures présentées vont de l'analyse cinématique des galaxies de l'échantillon spectroscopique MASSIV aux cartes de l'échantillon de simulations de fusion MIRAGE introduit dans le chapitre 2. Les articles auxquels j'ai participé au travers de la collaboration MASSIV sont eux présentés dans une deuxième annexe. Ces articles se focalisent sur l'analyse cinématique, et proposant une première classification de l'échantillon, ainsi qu'une estimation du taux de fusions dans l'Univers dans le domaine spectral couvert ($1.0 < z < 1.7$). Cette thèse a de plus été l'occasion de développer de plusieurs programmes informatiques. Dans une dernière annexe, je présente les trois codes les plus longs en termes de temps de développements. Ces codes sont à l'image de cette thèse : ils font le lien entre les observations et les simulations numériques de galaxies. Ces codes, qui ont fait l'objet d'un effort particulier sur l'écriture, sont mis à disposition de la communauté scientifique grâce au à l'hébergement de code en ligne. Ils ont vocation à être diffusés et améliorés au cours des prochaines années.

Cartes complémentaires



Résumé

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A.3 Champs de vitesses de pseudo-observations SINFONI [Chapitre 5]	239

Cette première annexe regroupe les différentes cartes qui n'ont pas pu être présentée dans le corps de la thèse du fait de leur volume trop important. Chaque partie de cette annexe fait référence à un chapitre particulier, qui est indiquée dans le titre de la section.

A.1 Kinemetry sur les cartes MASSIV [Chapitre 1]

La décomposition en harmoniques sphériques de chacune de galaxies MASSIV est présentée dans cette partie des annexes. Les paramètres géométriques des ellipses le long duquel l'information cinématique est extraite ont été obtenu grâce à la routine `KINEGEN`. Pour chaque galaxie ayant suffisamment de signal détecté (un minimum de 5 pixels détecté au dessus de 3σ pour la raie d'émission $H\alpha$) de l'échantillon MASSIV, la décomposition du champ de vitesses et du champ de dispersion de vitesses est montrée, en illustrant notamment l'expansion harmonique sur une ellipse.

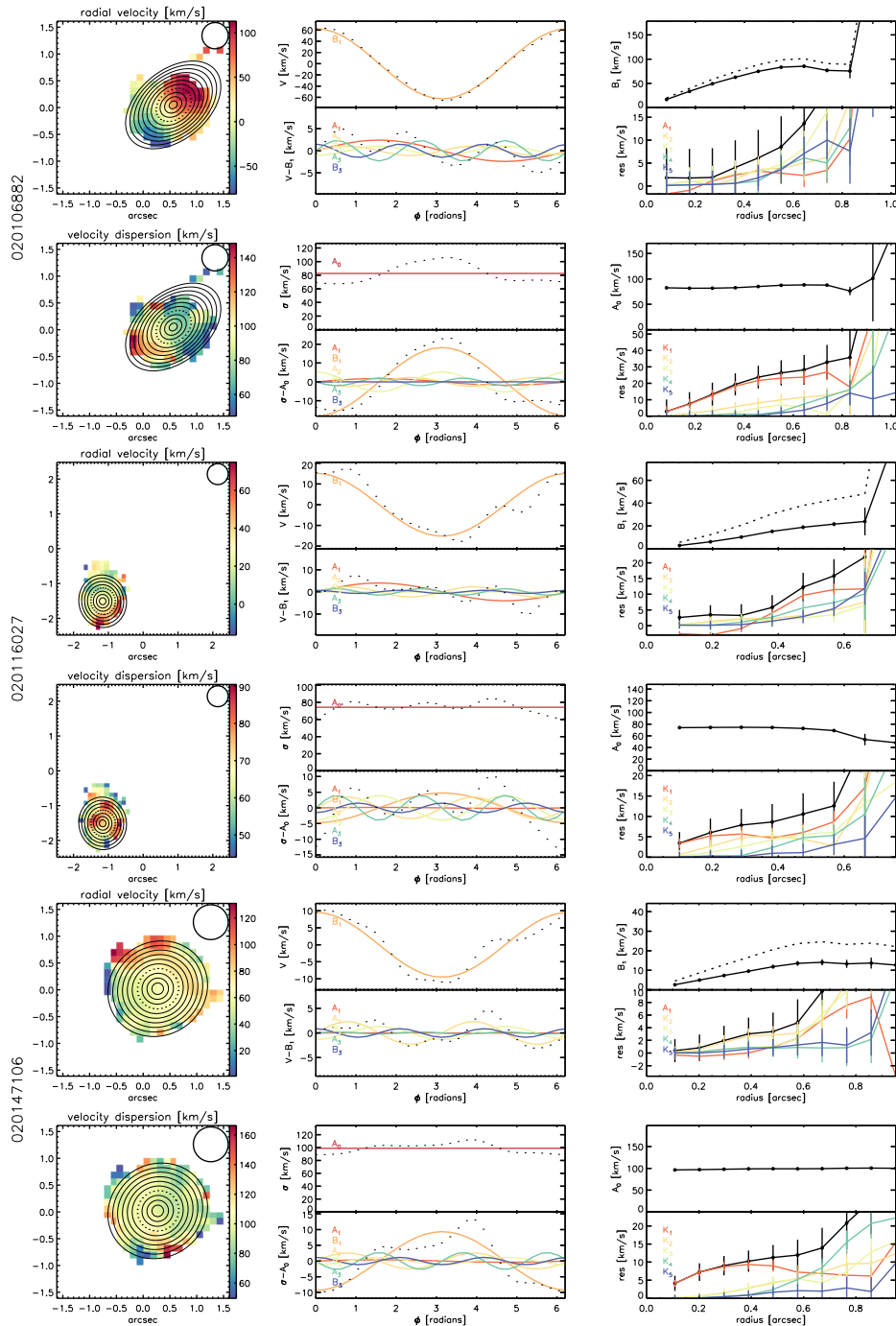
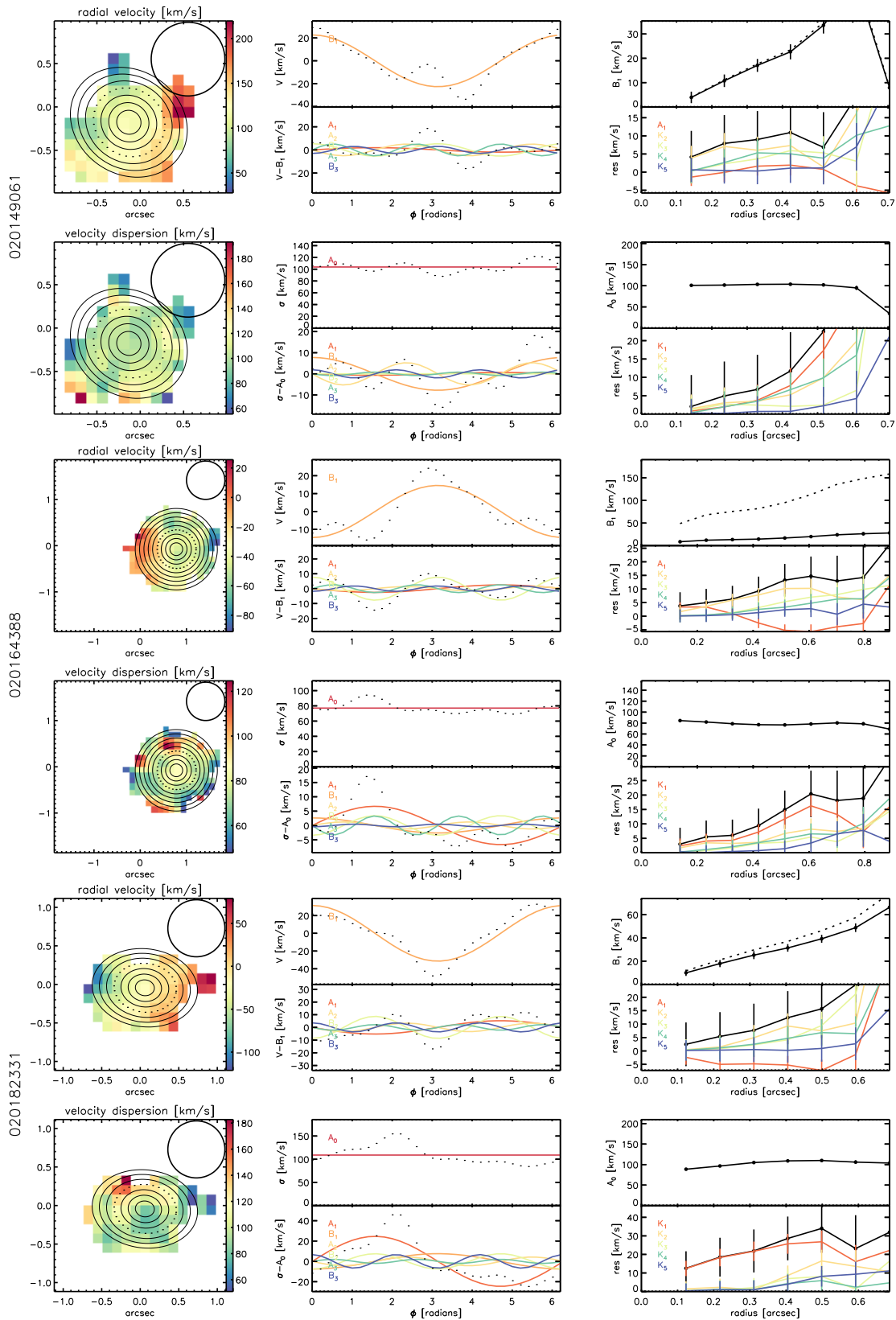
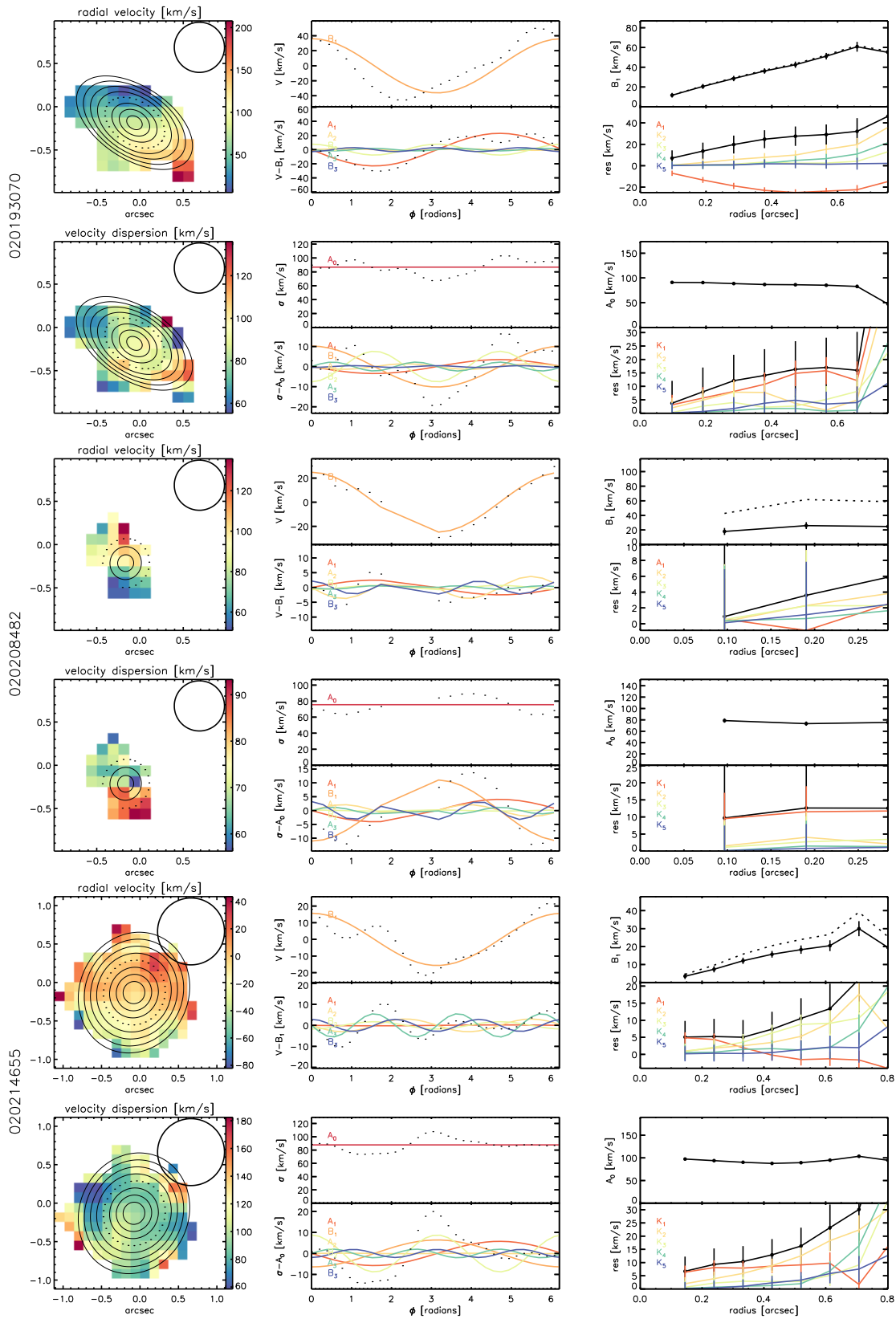


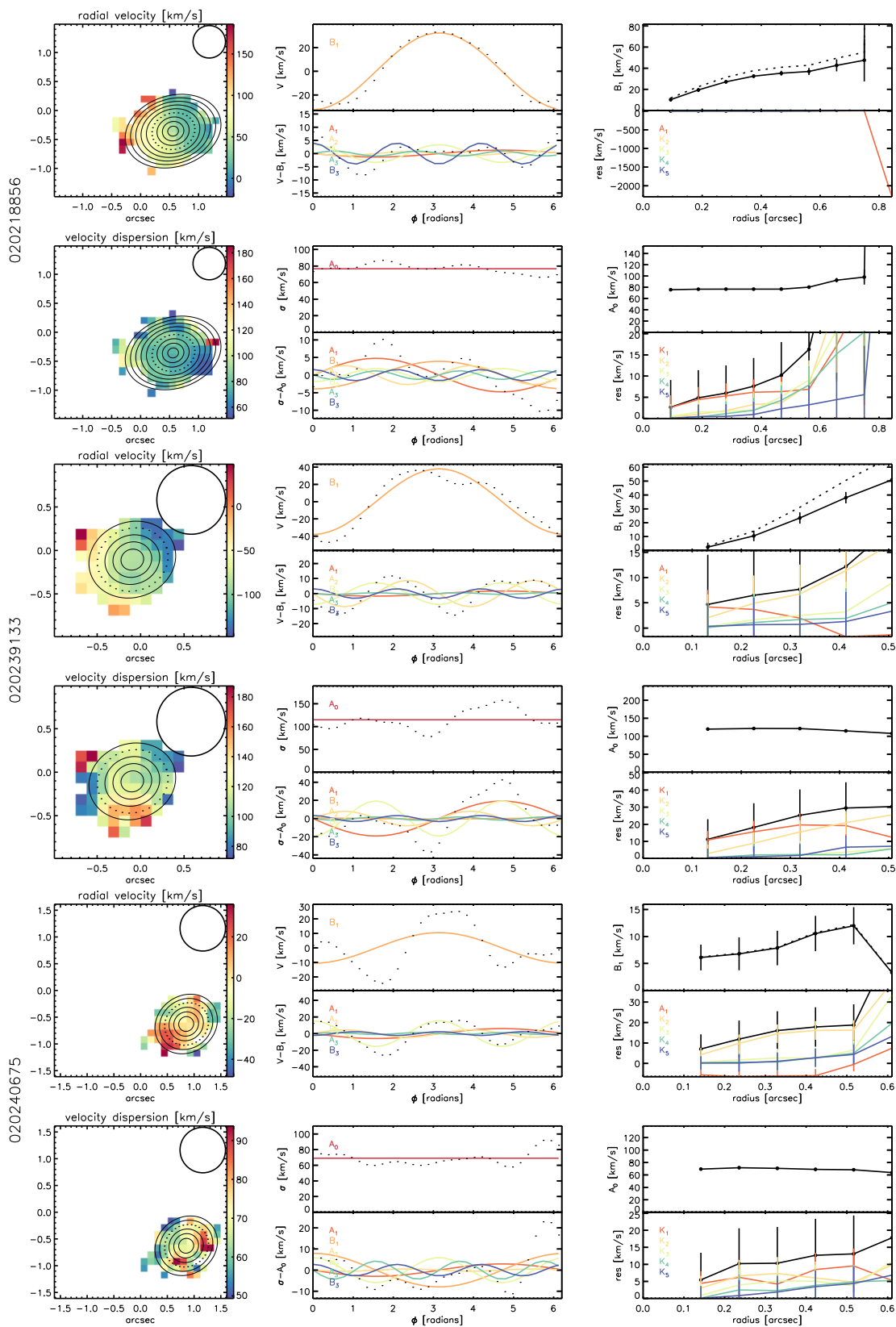
Figure A.1 – Kinemetry pour trois galaxies de l'échantillon MASSIV. Pour chaque carte des moments (vitesses radiales et dispersion de vitesses) : À gauche : carte des moments avec les ellipses superposées. Le cercle en haut à droite donne la taille angulaire de la PSF. Au milieu : Coefficients de l'expansion harmonique en fonction de l'angle d'azimut le long de l'ellipse tracée par des points noirs. Le panneau du haut montre la mesure du moment (points noirs) le long de l'ellipse, tandis que le panneau du bas montre les résidus une fois le moment principal ajusté (B_1 pour la vitesses radiales, A_0 pour la dispersion de vitesses) soustrait (points noirs). Les harmoniques d'ordre supérieur ajustées sont aussi tracées. À droite : Le panneau du haut montre la dépendance radiale du coefficient harmonique principal (courbe noire). Dans le cas de l'expansion sur le champs de vitesses, le terme B_1 est aussi déprojeté de l'inclinaison d'entrée (courbe en pointillés). Le panneau du bas montre la dépendance radiale des coefficients résiduels. La courbe noire montre la somme quadratique jusqu'à $n=5$ des coefficients résiduels. L'identifiant de la galaxie MASSIV est indiqué sur la gauche de chaque carte.



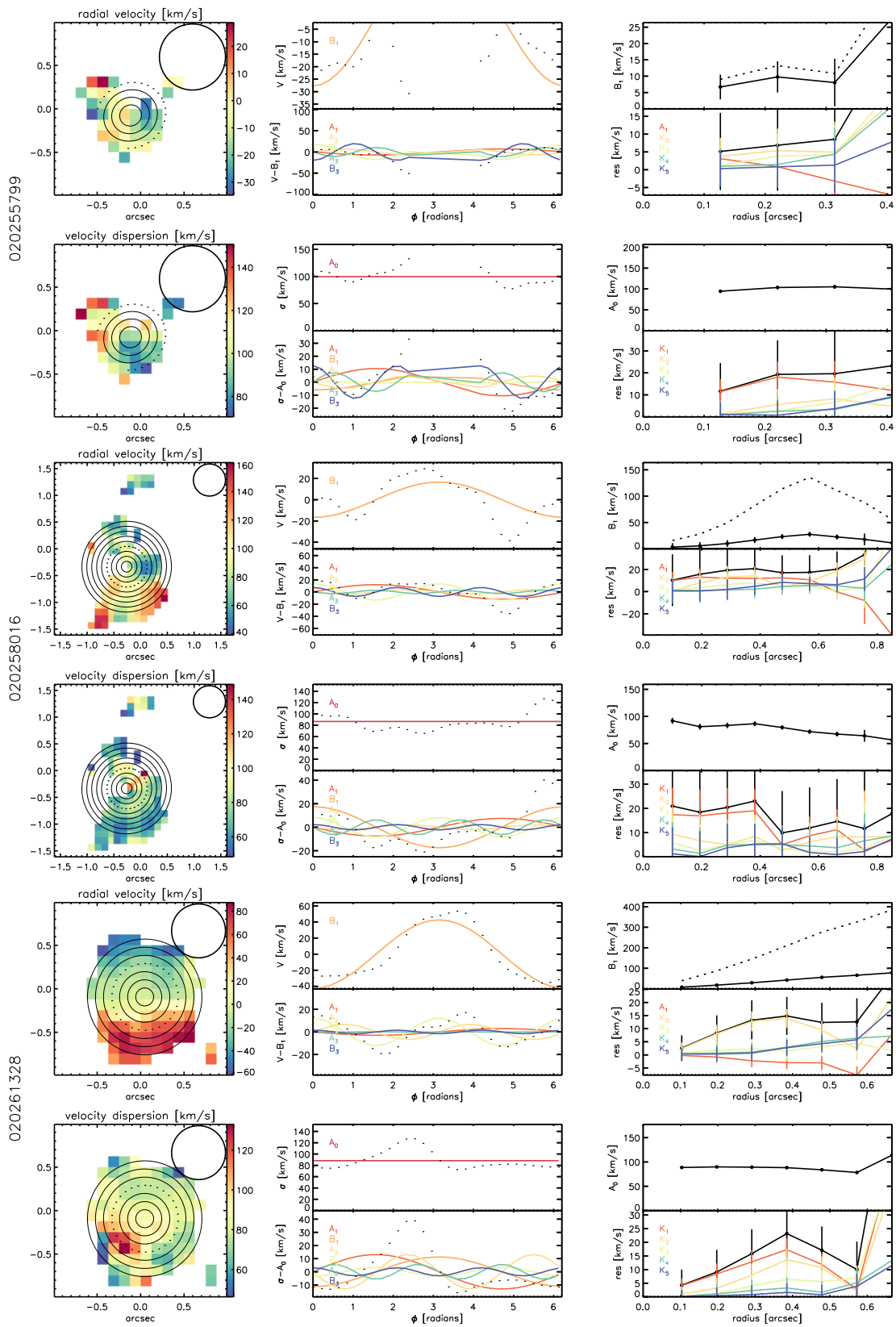
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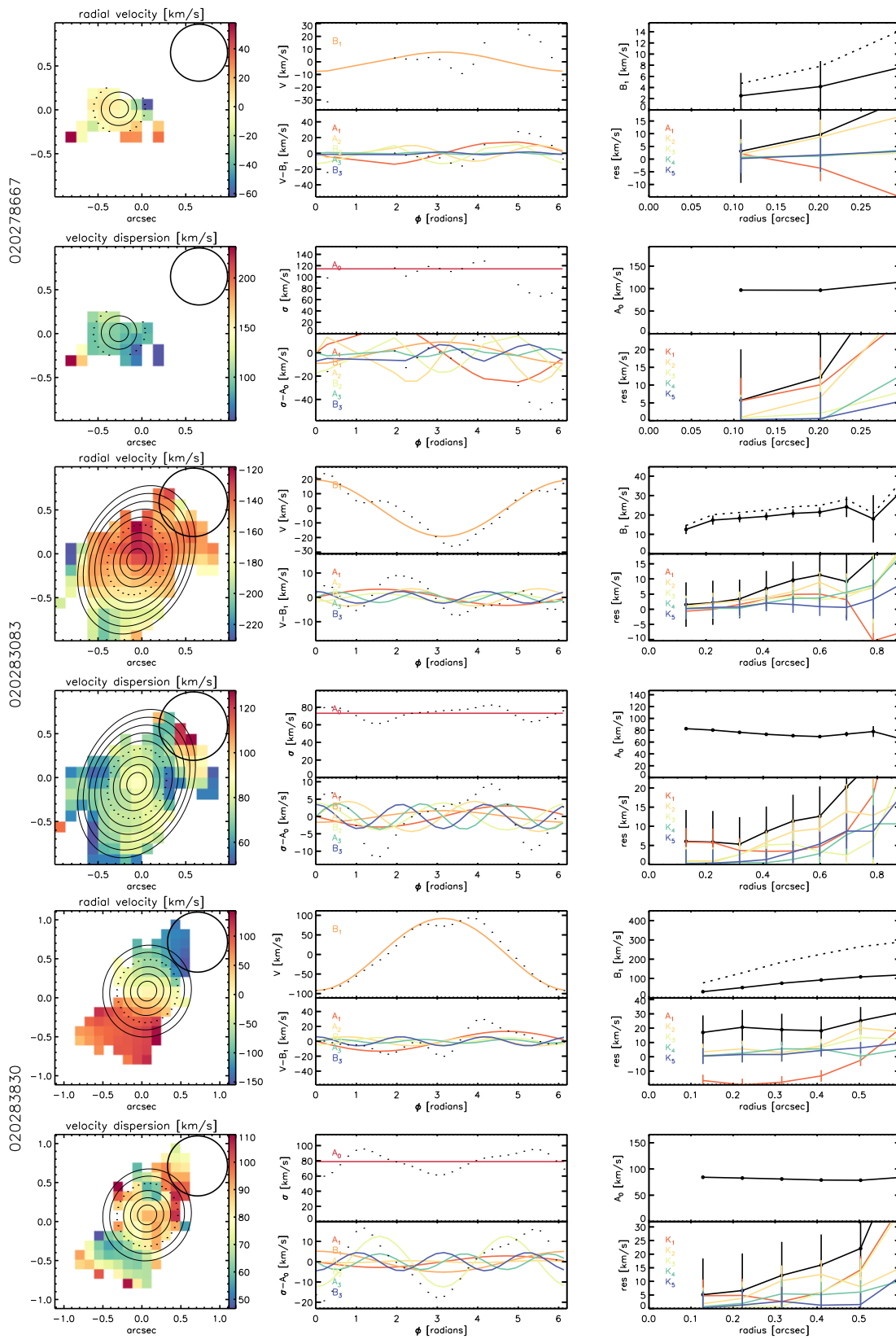
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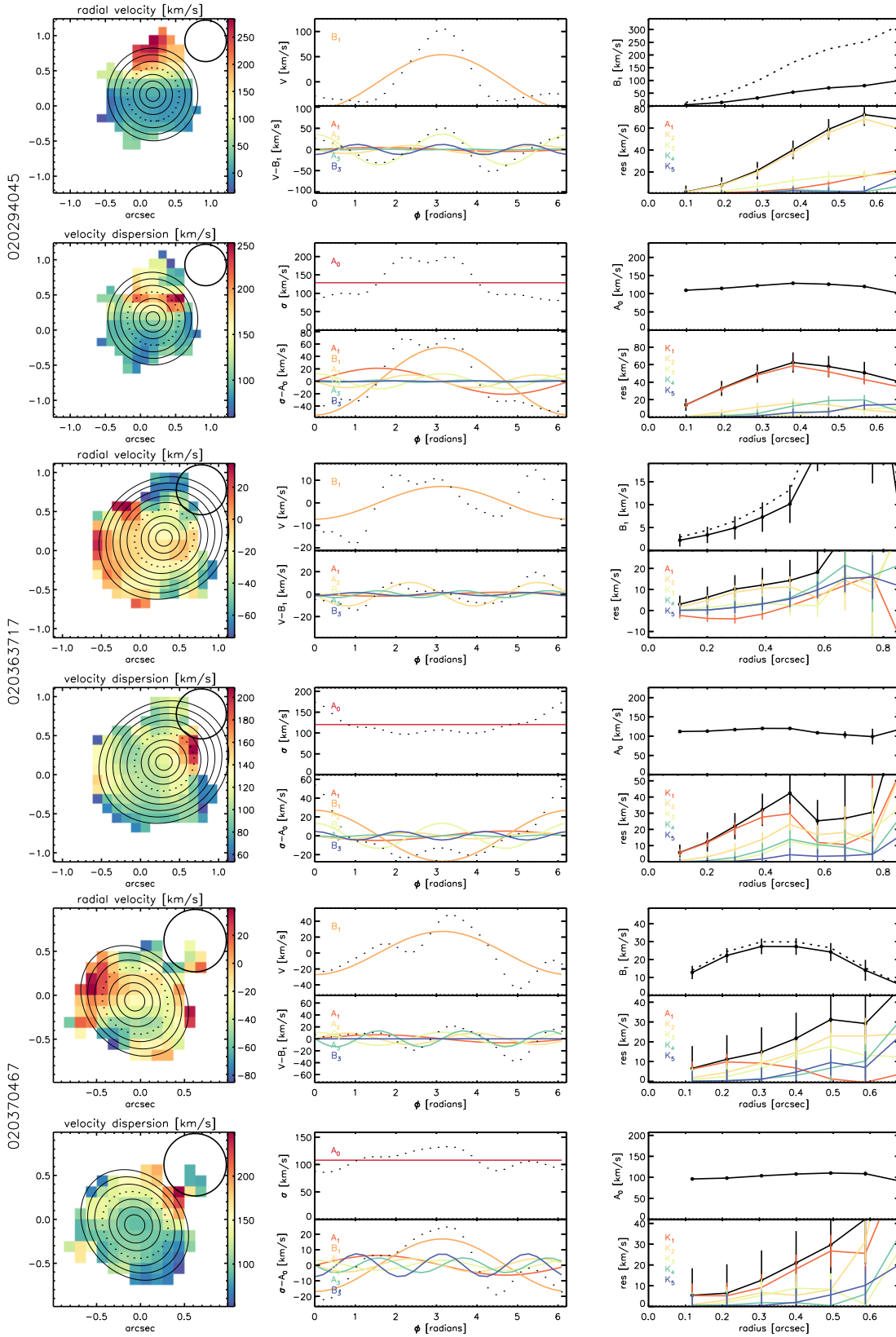
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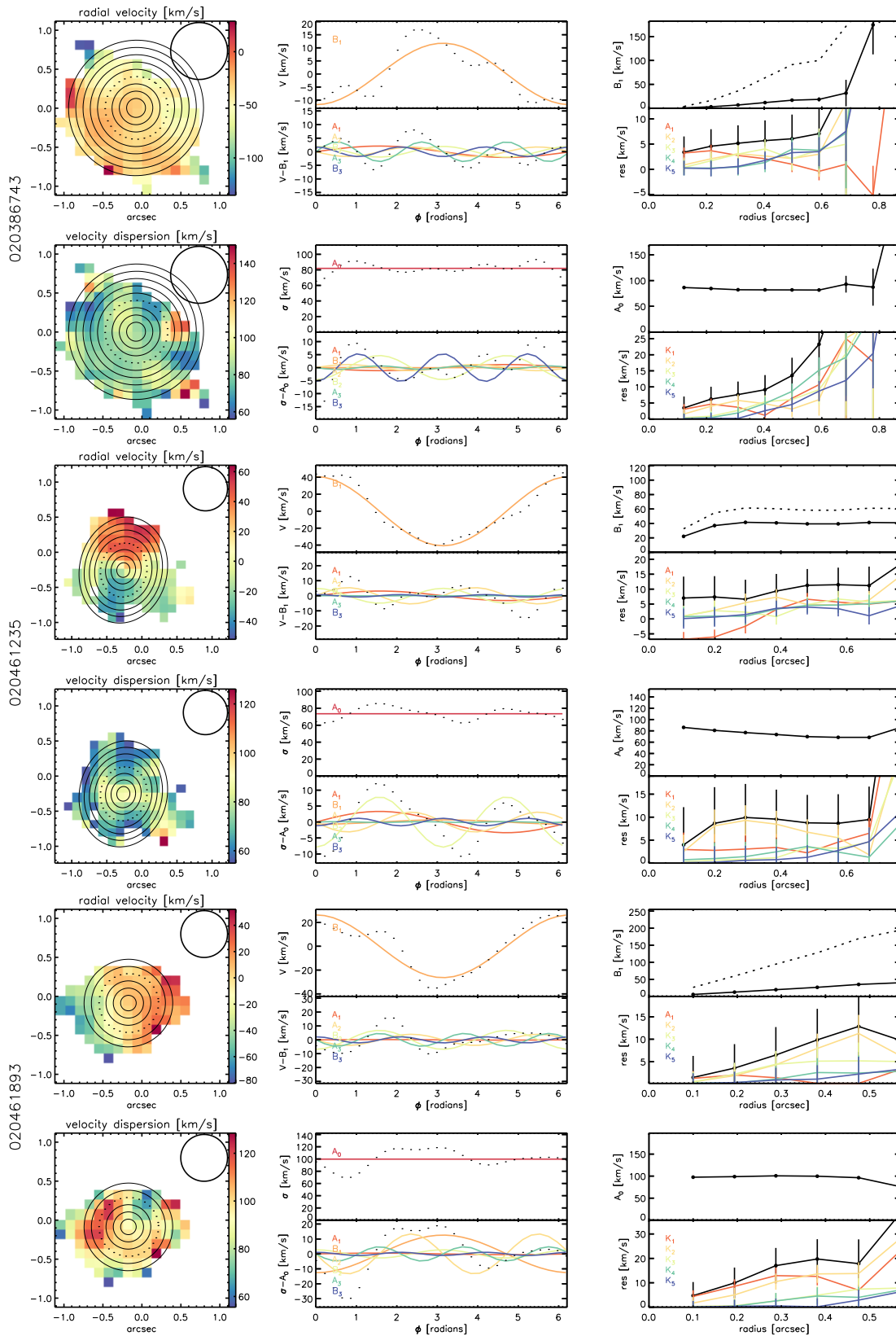
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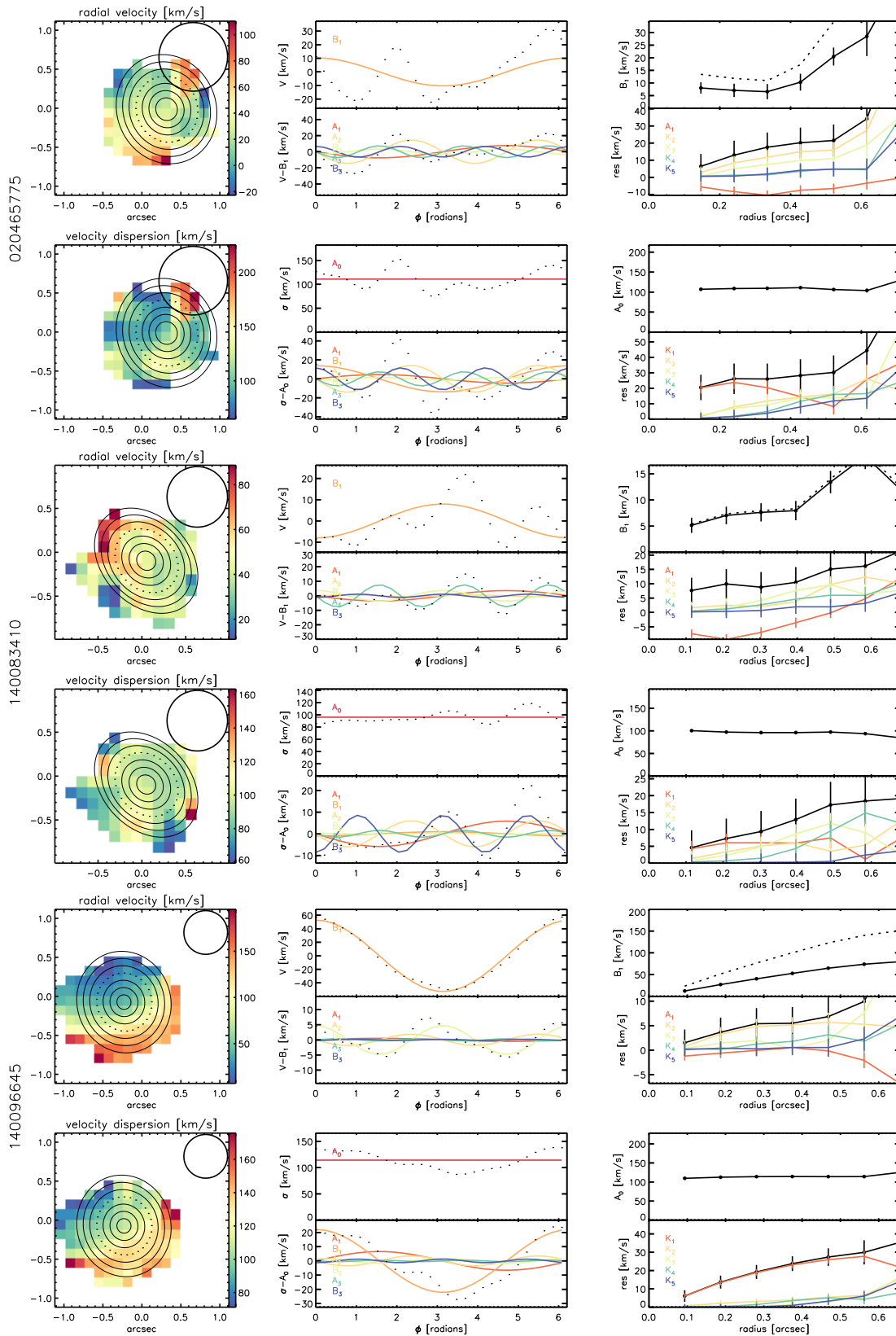
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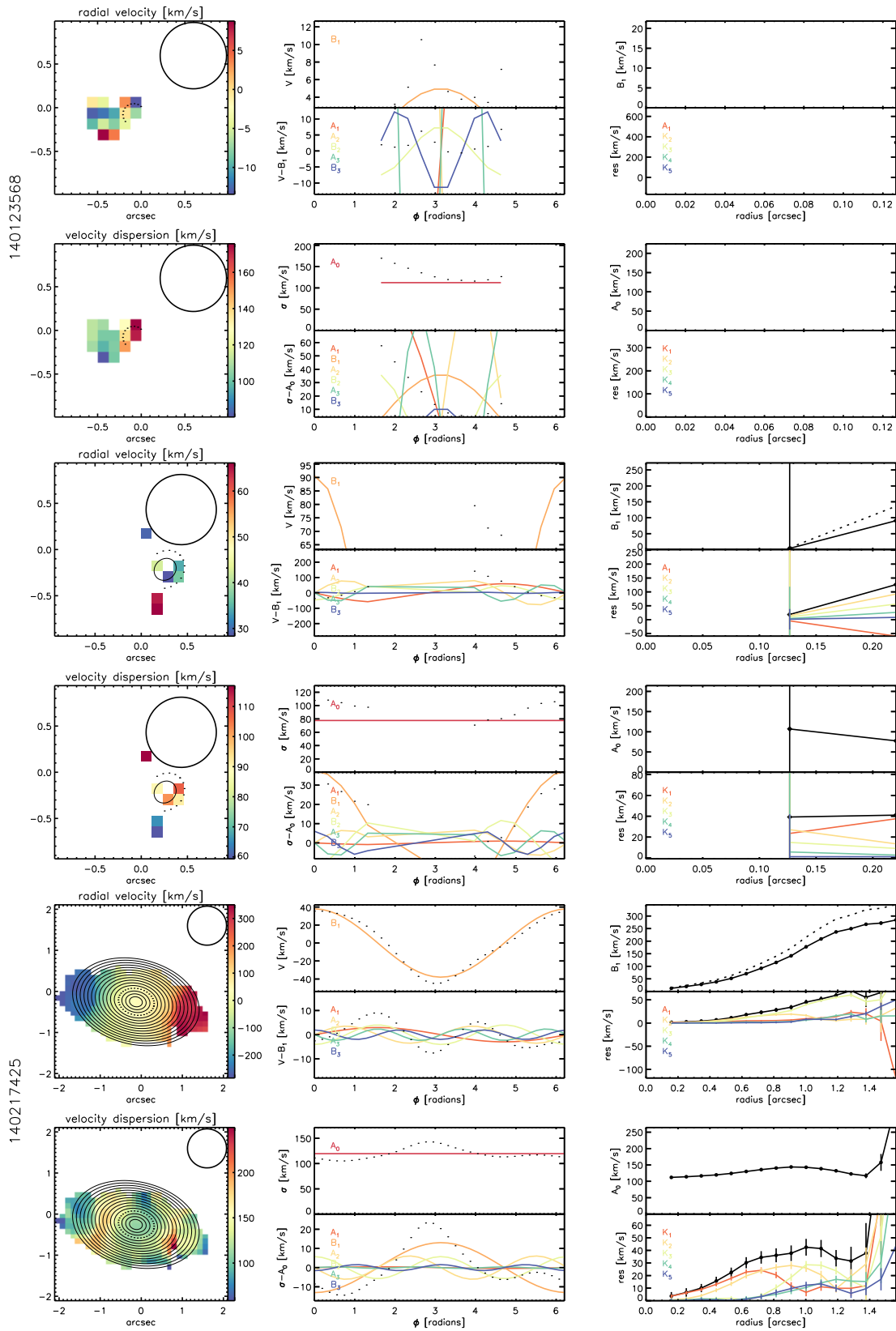
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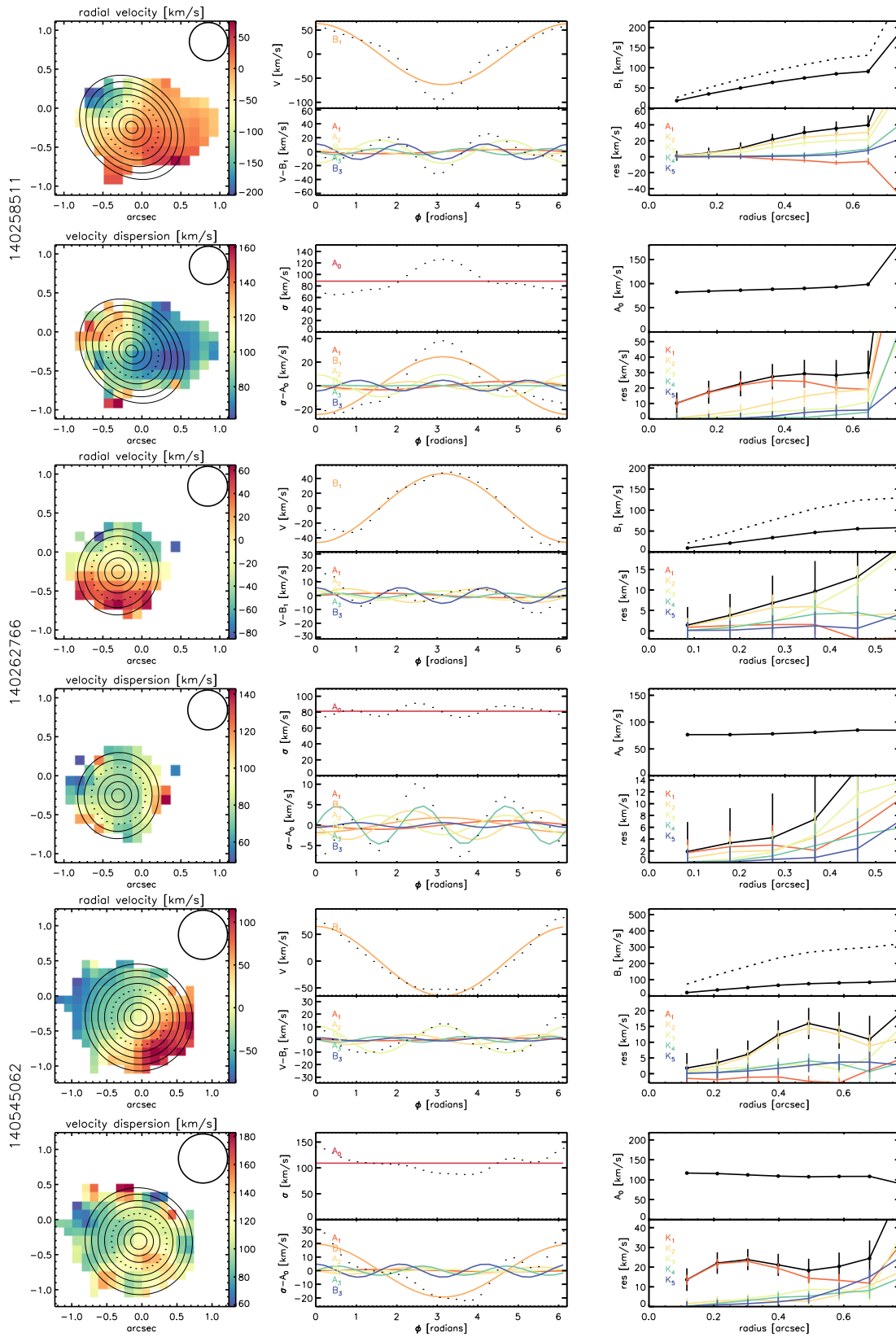
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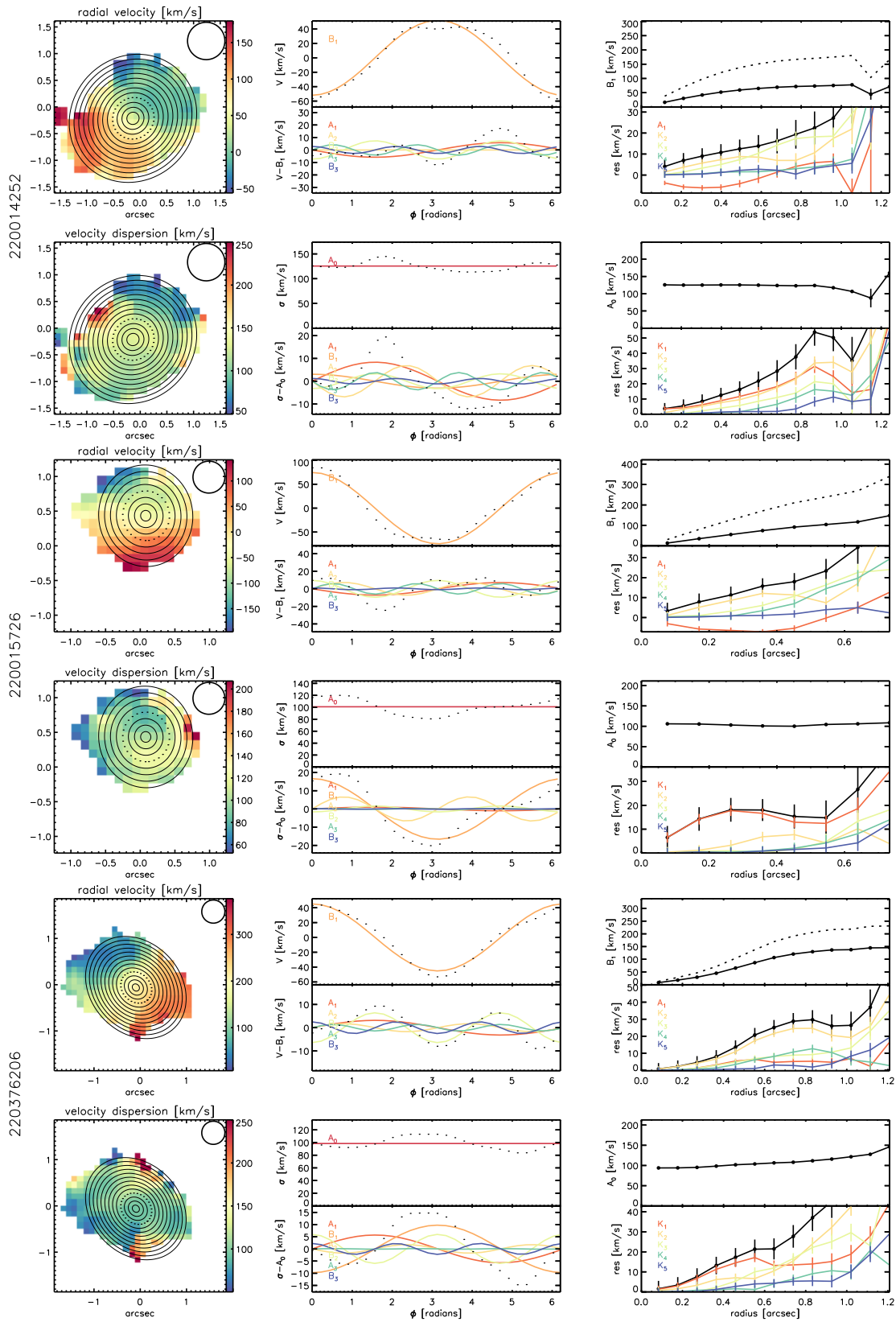
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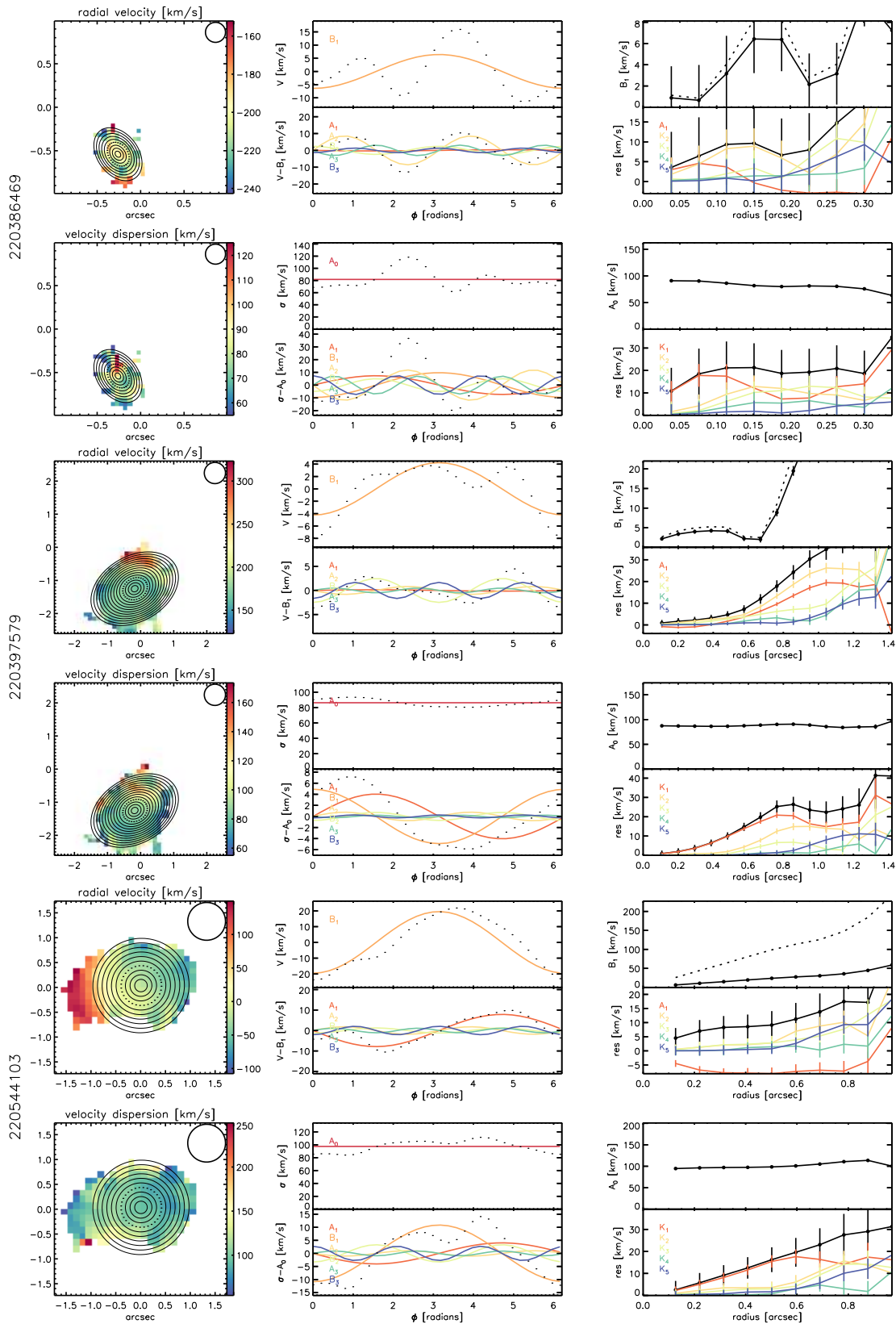
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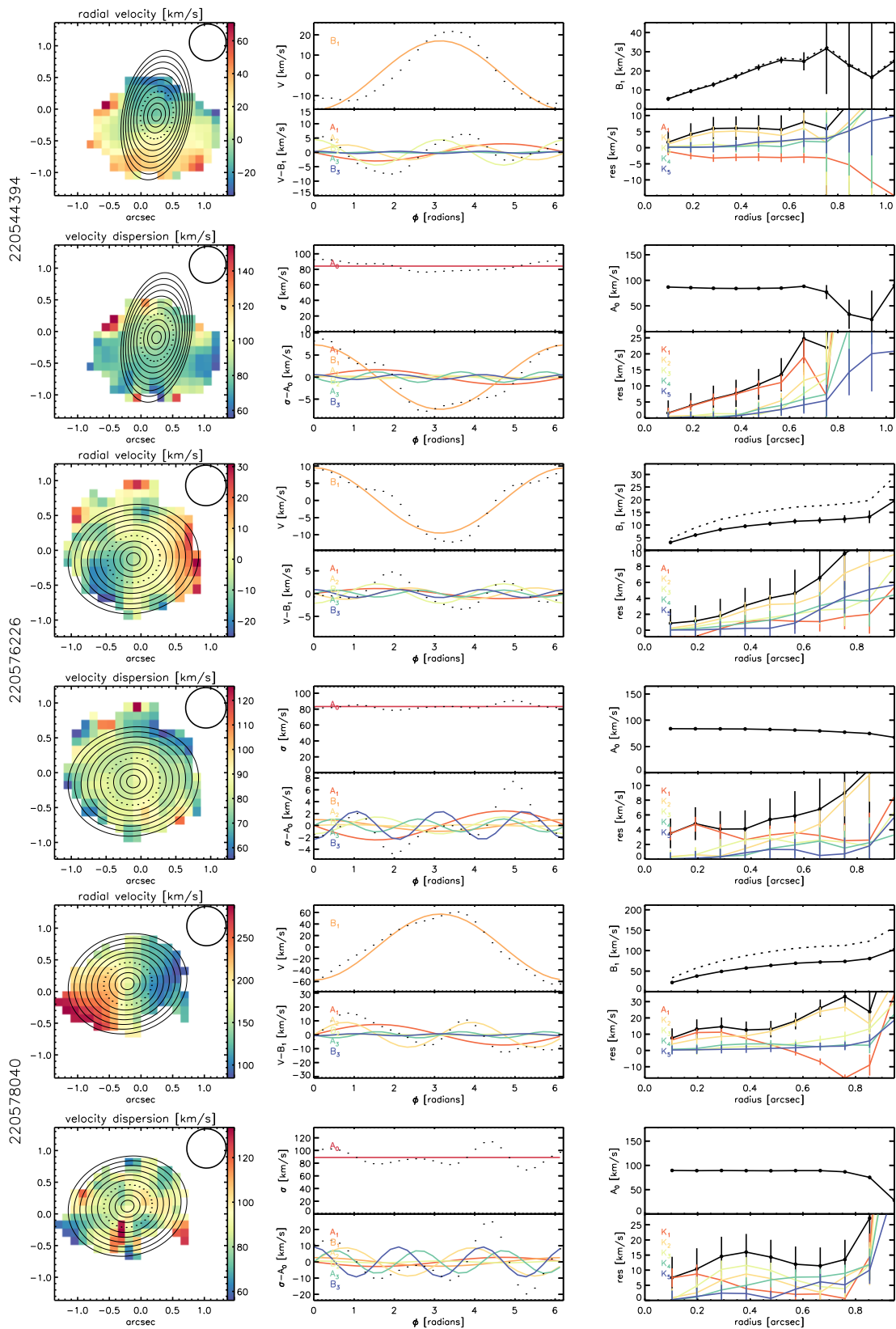
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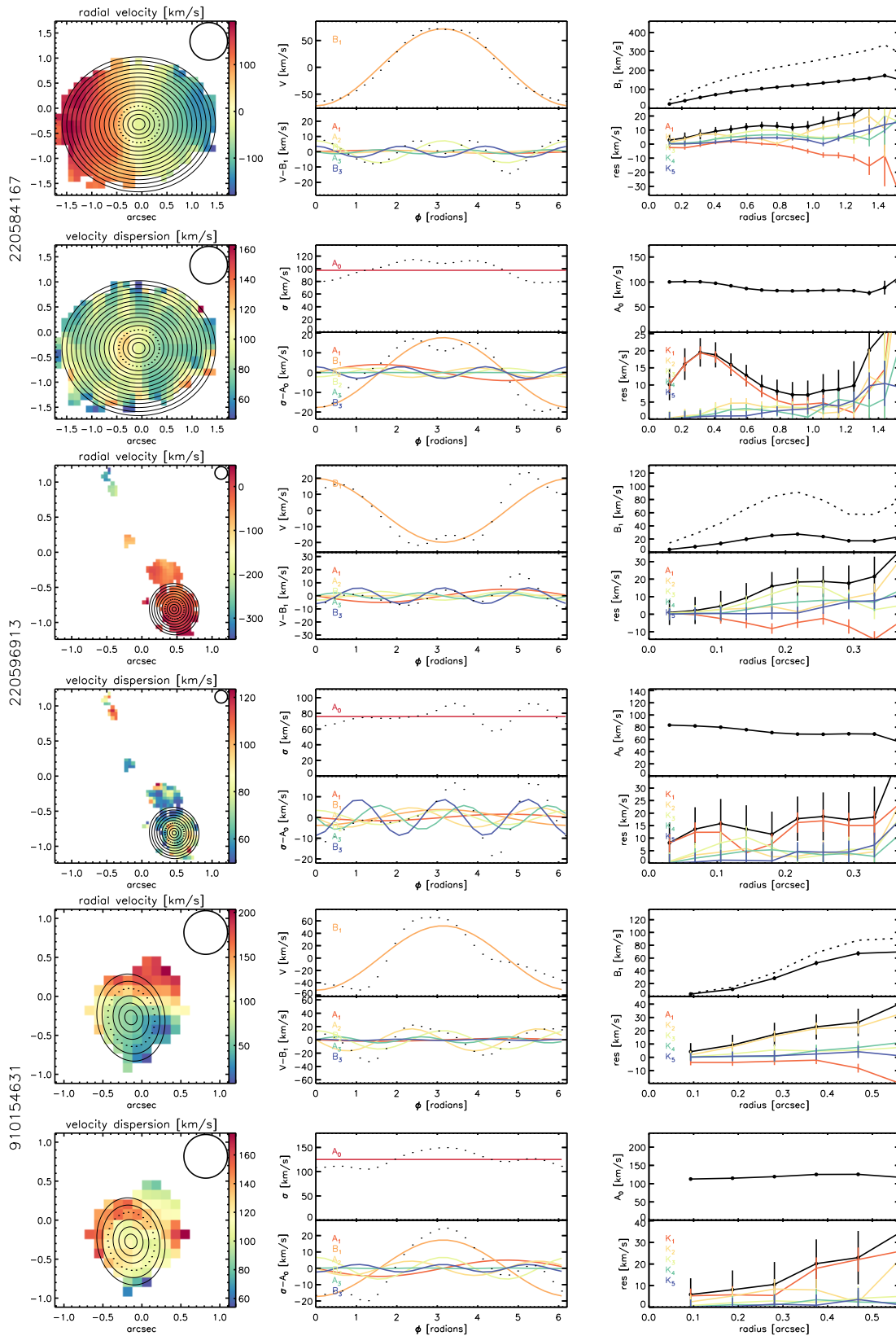
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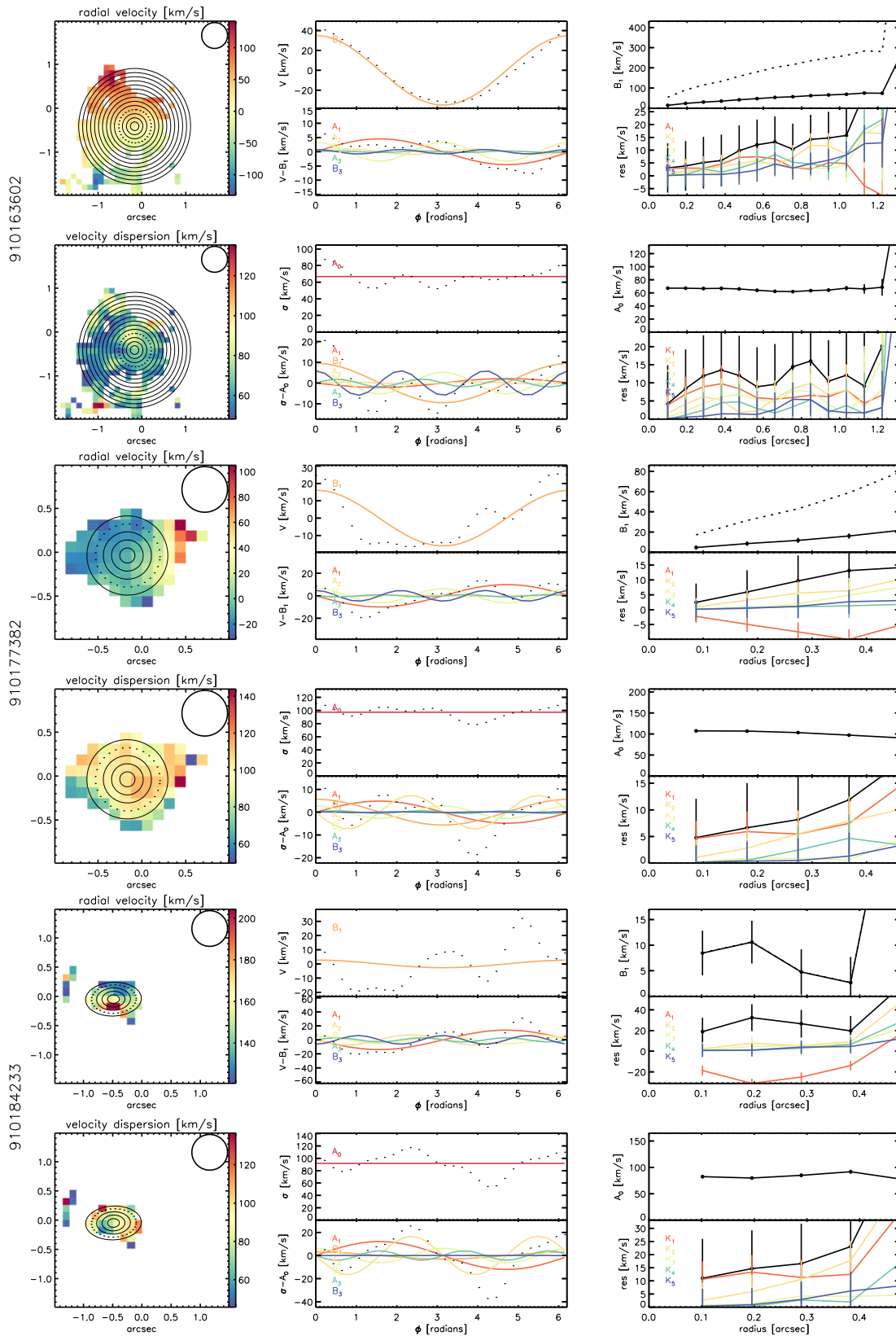
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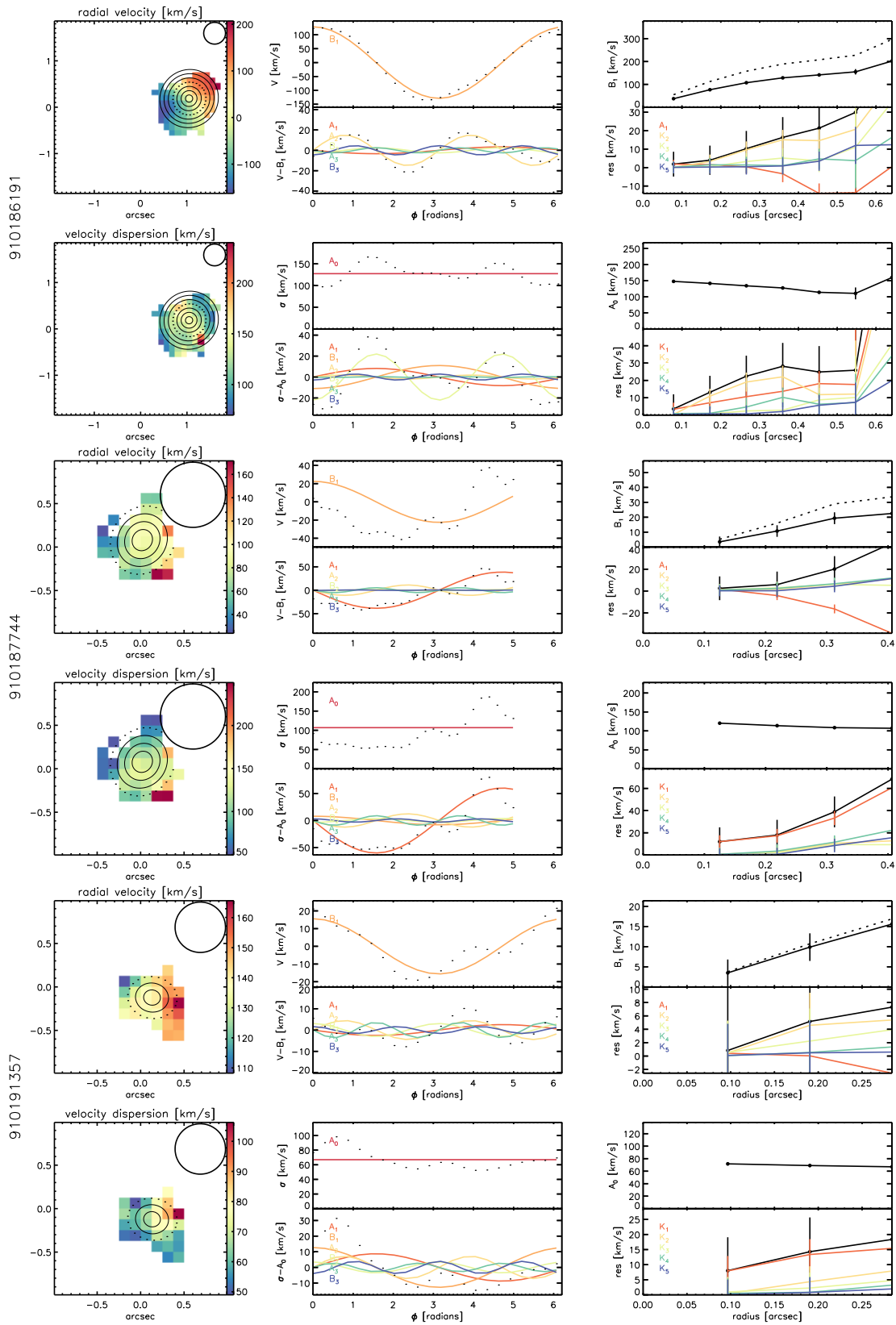
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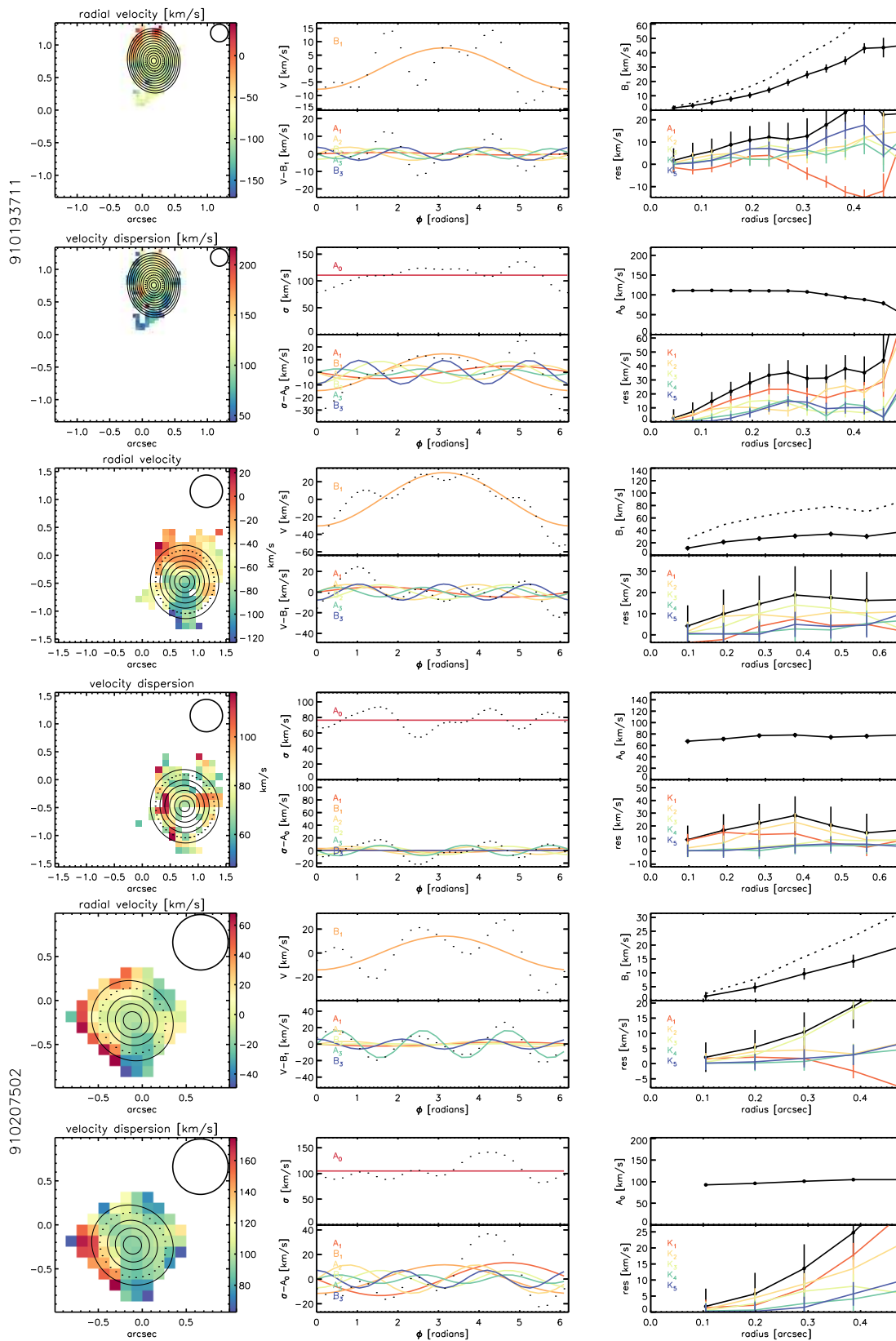
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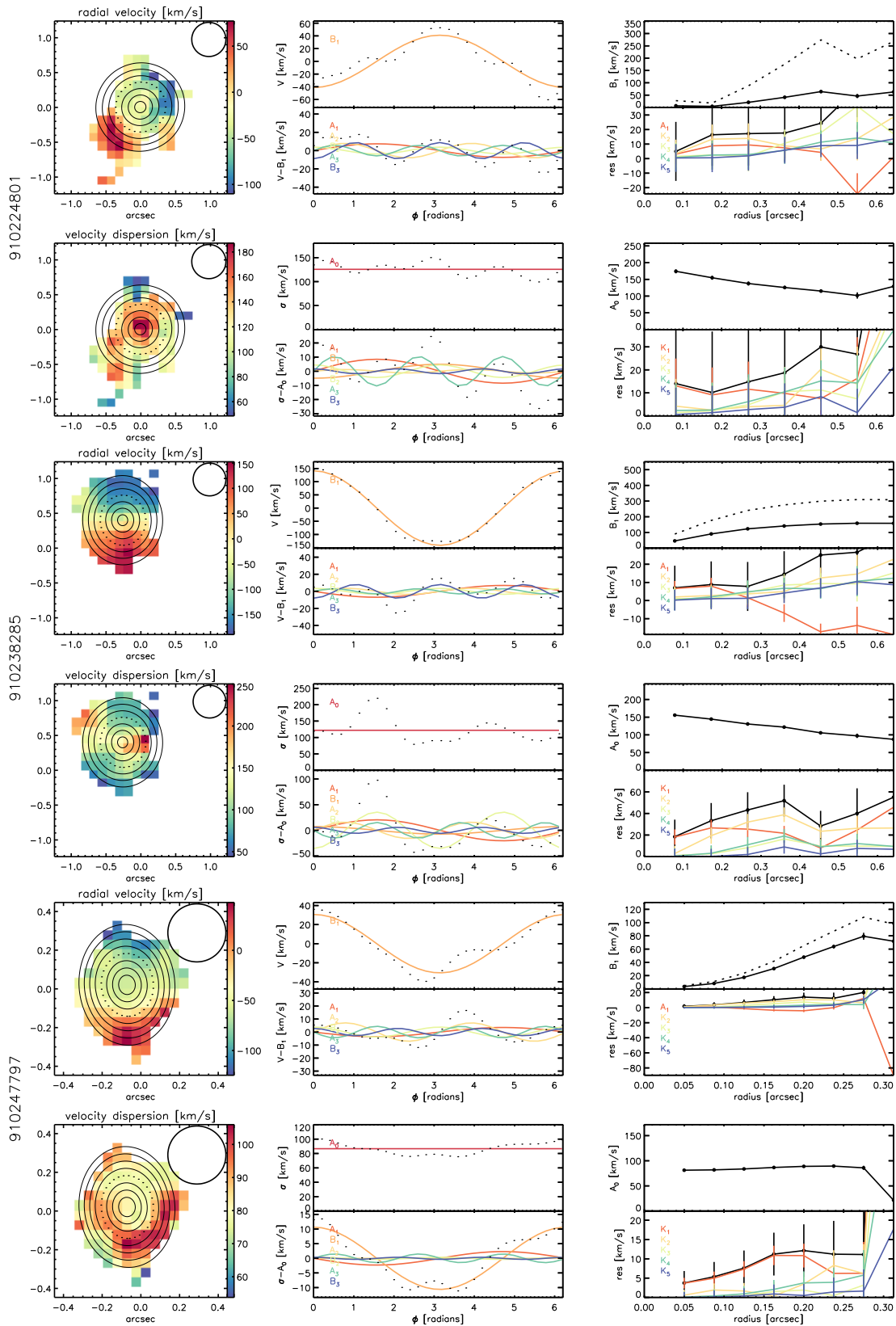
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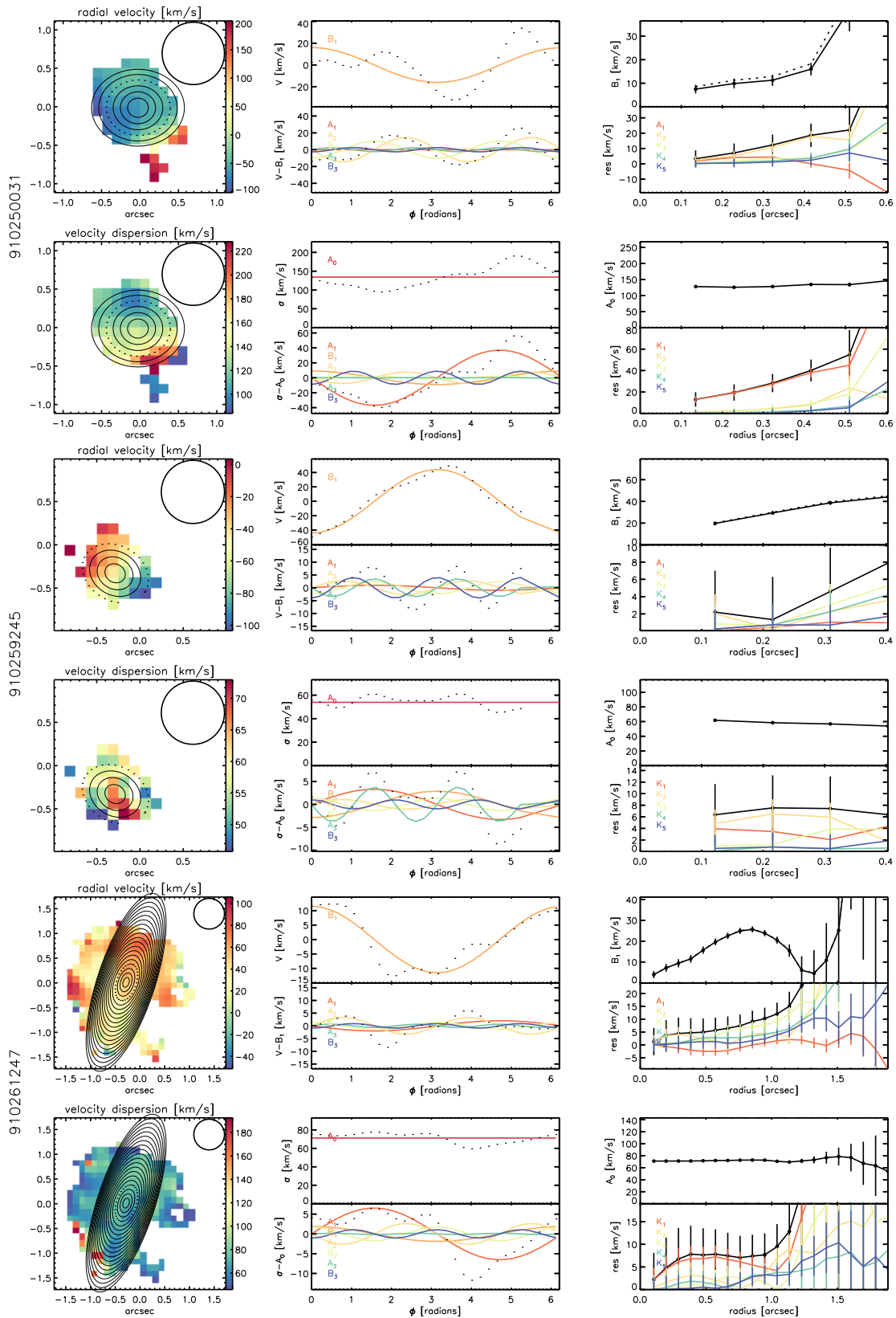
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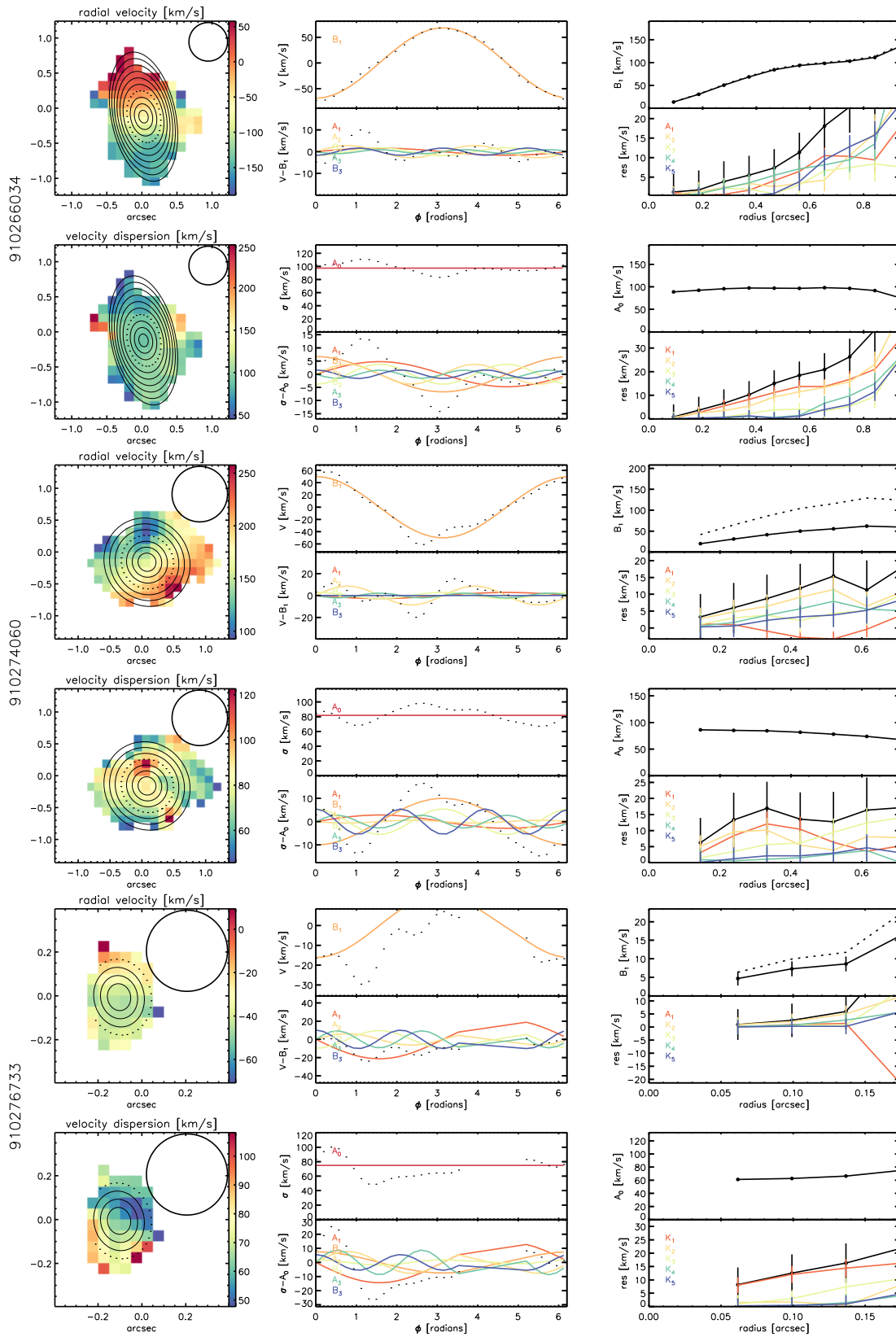
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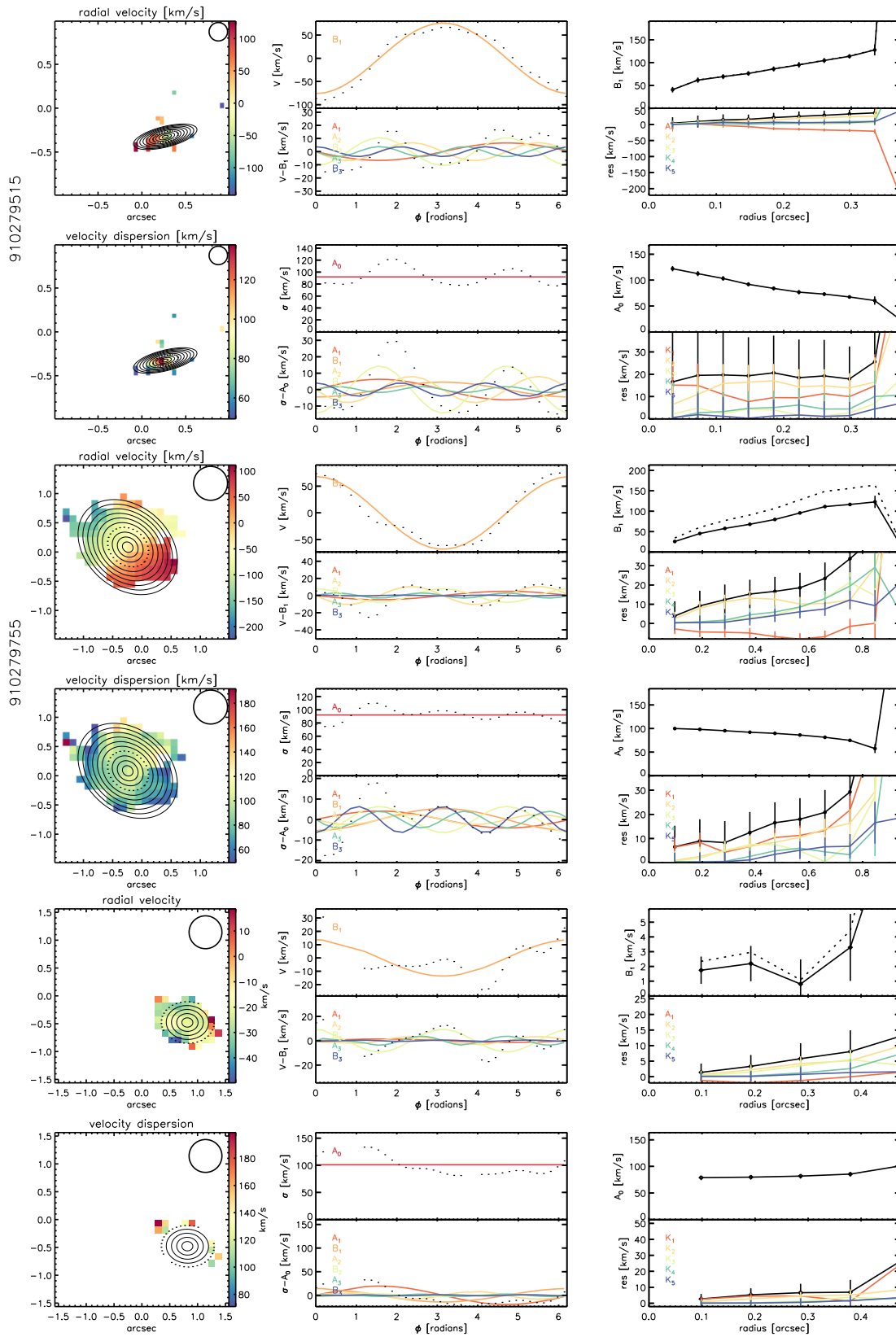
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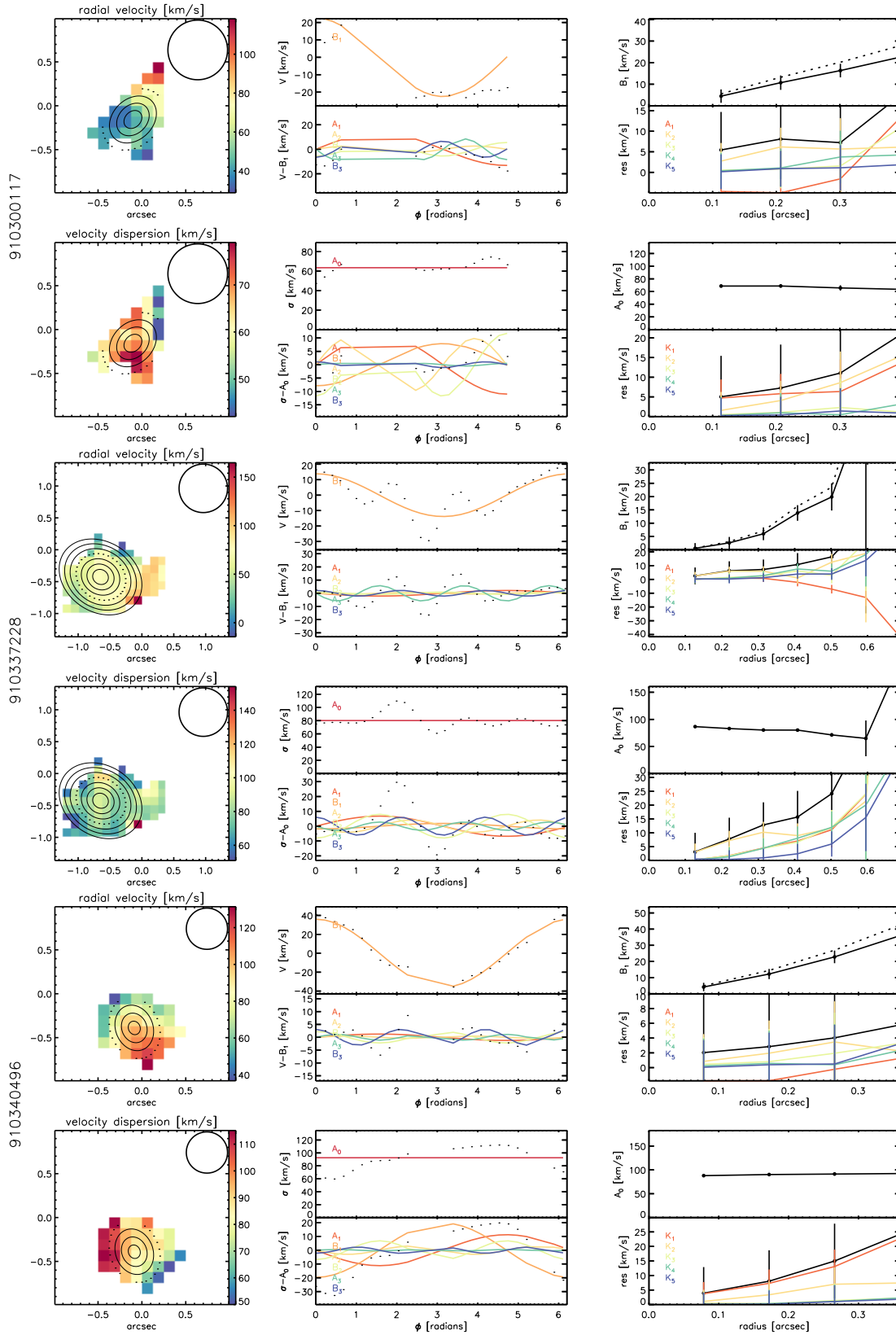
Suite de la figure A.1.



Suite de la figure A.1.



Suite de la figure A.1.



Suite de la figure A.1.

A.2 Cartes du gaz et des étoiles des simulations MIRAGE [Chapitre 3]

Dans le chapitre 3, je présente l'échantillon de simulations idéalisées de fusions de galaxies jeunes MIRAGE, qui est l'un des produits majeur de cette thèse. Cet échantillon explore les paramètres initiaux d'orientation et de masse des deux disques progéniteurs, pour une énergie orbitale spécifique constante. Il est composé de 23 simulations explorant 800 Myr d'évolution : trois simulations de disques en évolution séculaire, et 20 simulations de fusions. Cette section de l'annexe présente les cartes de quantités hydrodynamiques et de flux stellaire réalisées avec le programme Aurora (5), et permet de mieux appréhender la dynamique et la morphologie de chacun des modèles numériques.

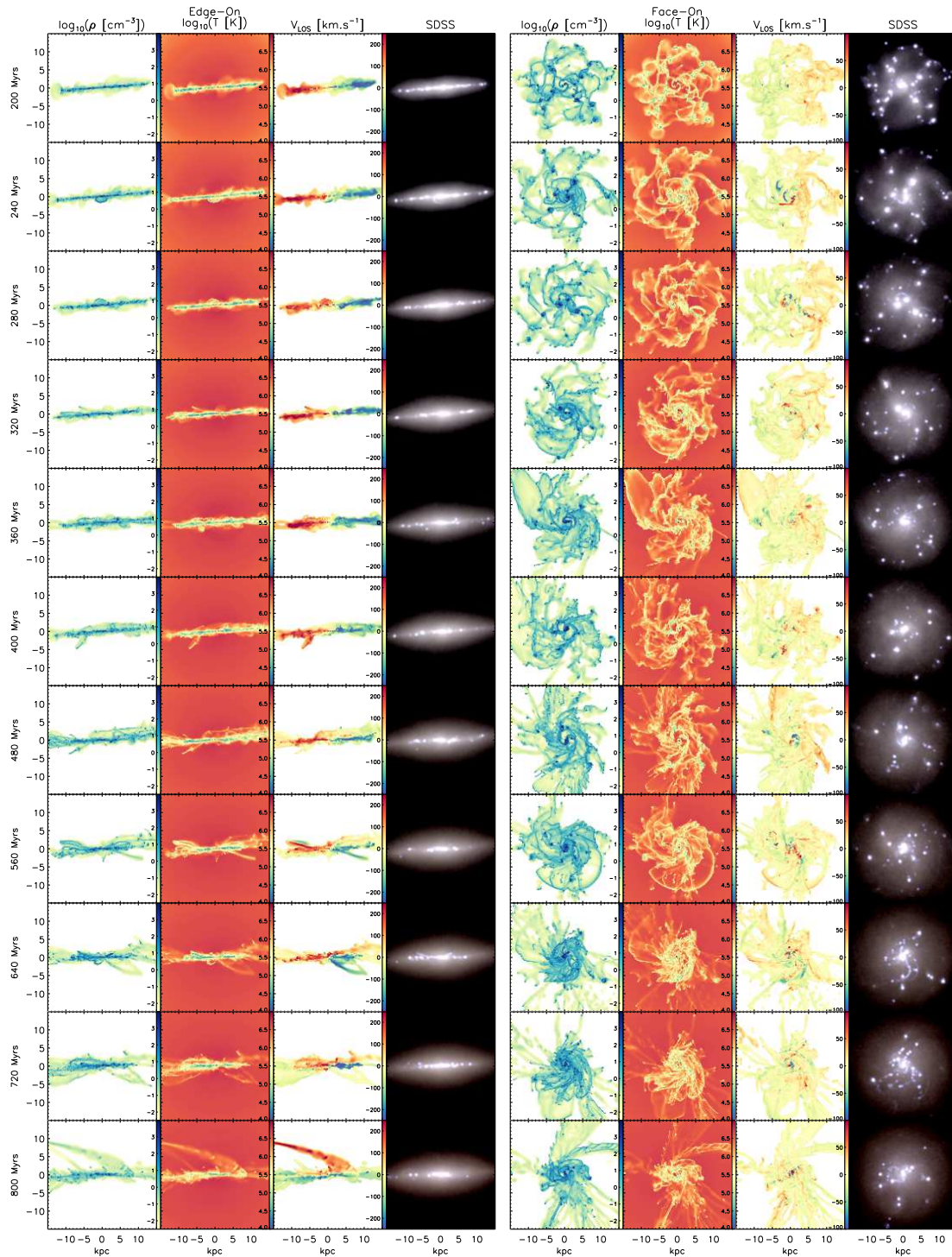
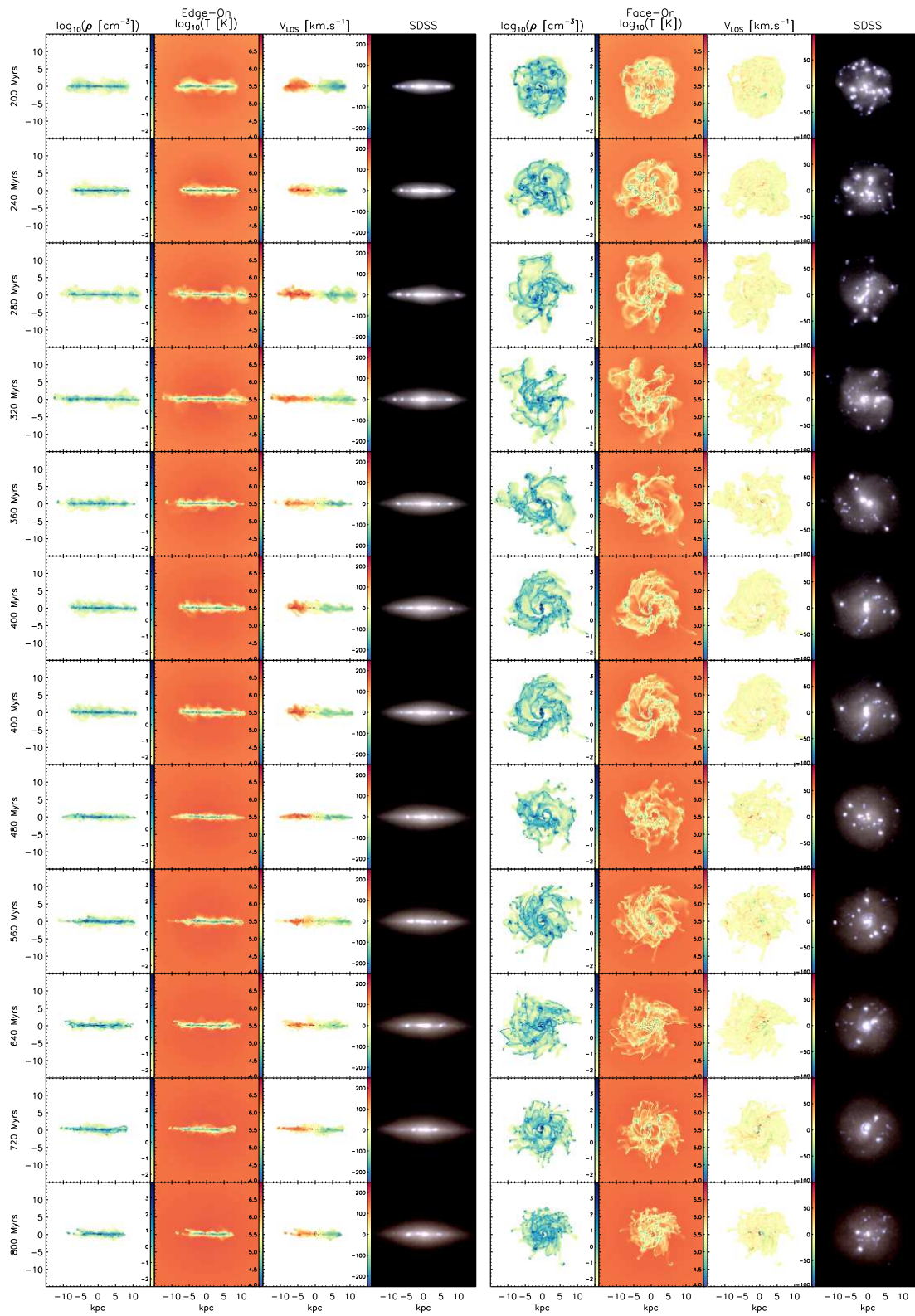
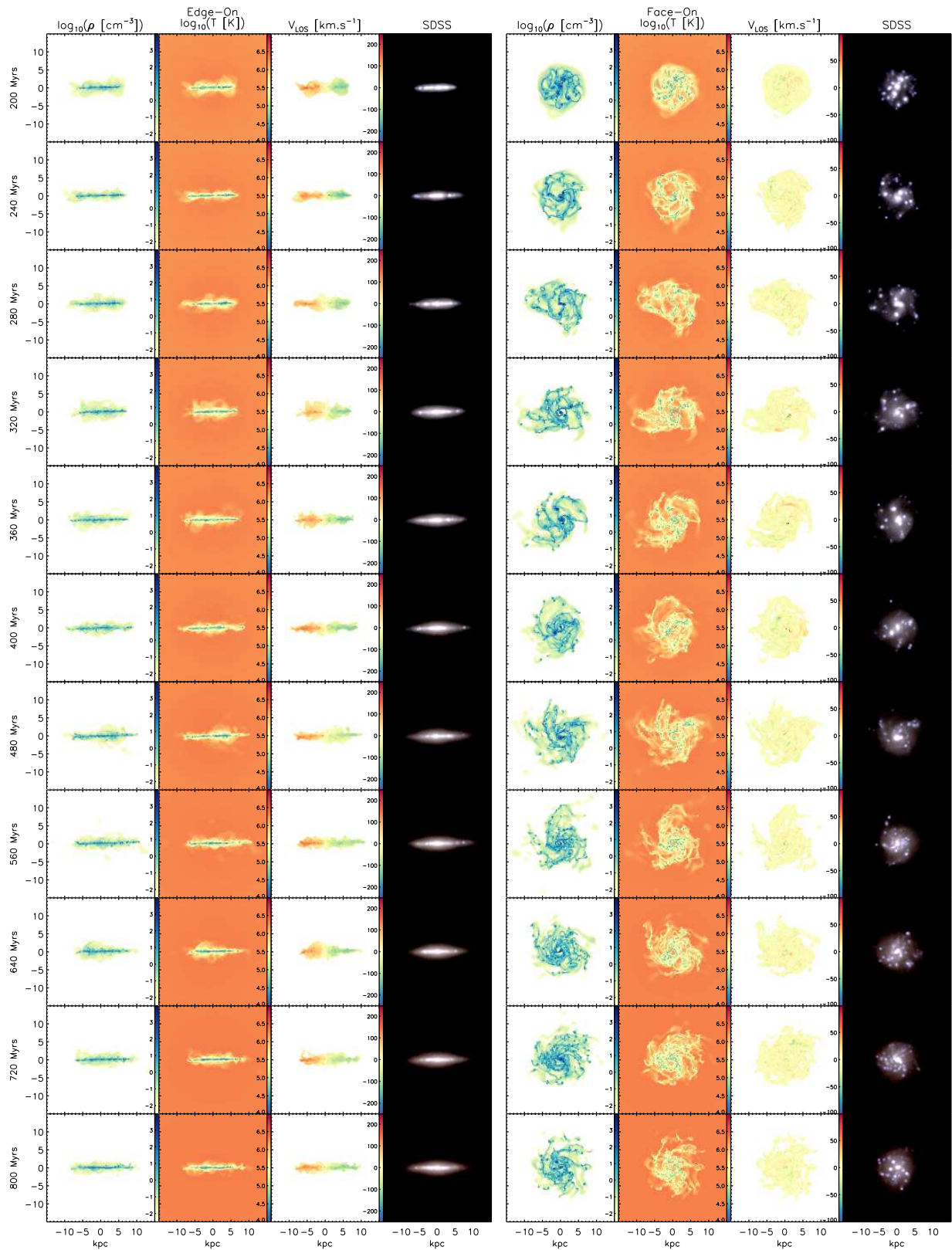


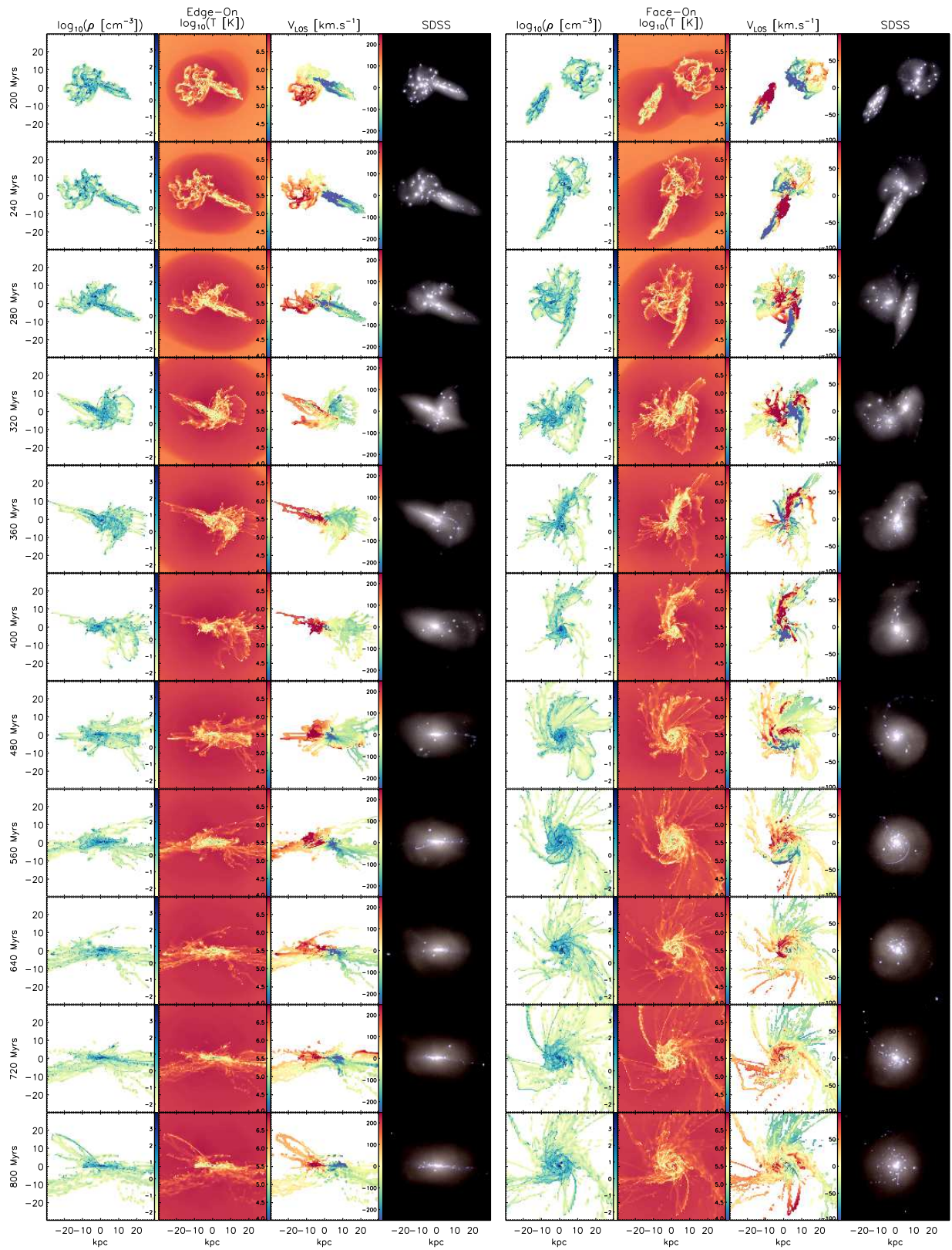
Figure A.2 – Projections orthographiques de la simulation G1. Deux lignes de visée orthogonales sont utilisées dans deux jeux de colonnes. Une projection du disque vu de face est présentée dans les quatre colonnes de gauche, et une projection du disque vu par la tranche est présentée dans les quatre colonnes de droite. Pour chaque projection, on affiche de gauche à droite la densité de gaz, la température du gaz, la vitesse Doppler du gaz, et une image composite réalisée à partir de trois pseudo-observations du continu stellaire simulant l'instrument SDSS dans les bandes ugr . Les cartes des quantités hydrodynamiques sont toutes calculées à partir de moyennes pondérées par la masse des cellules AMR.



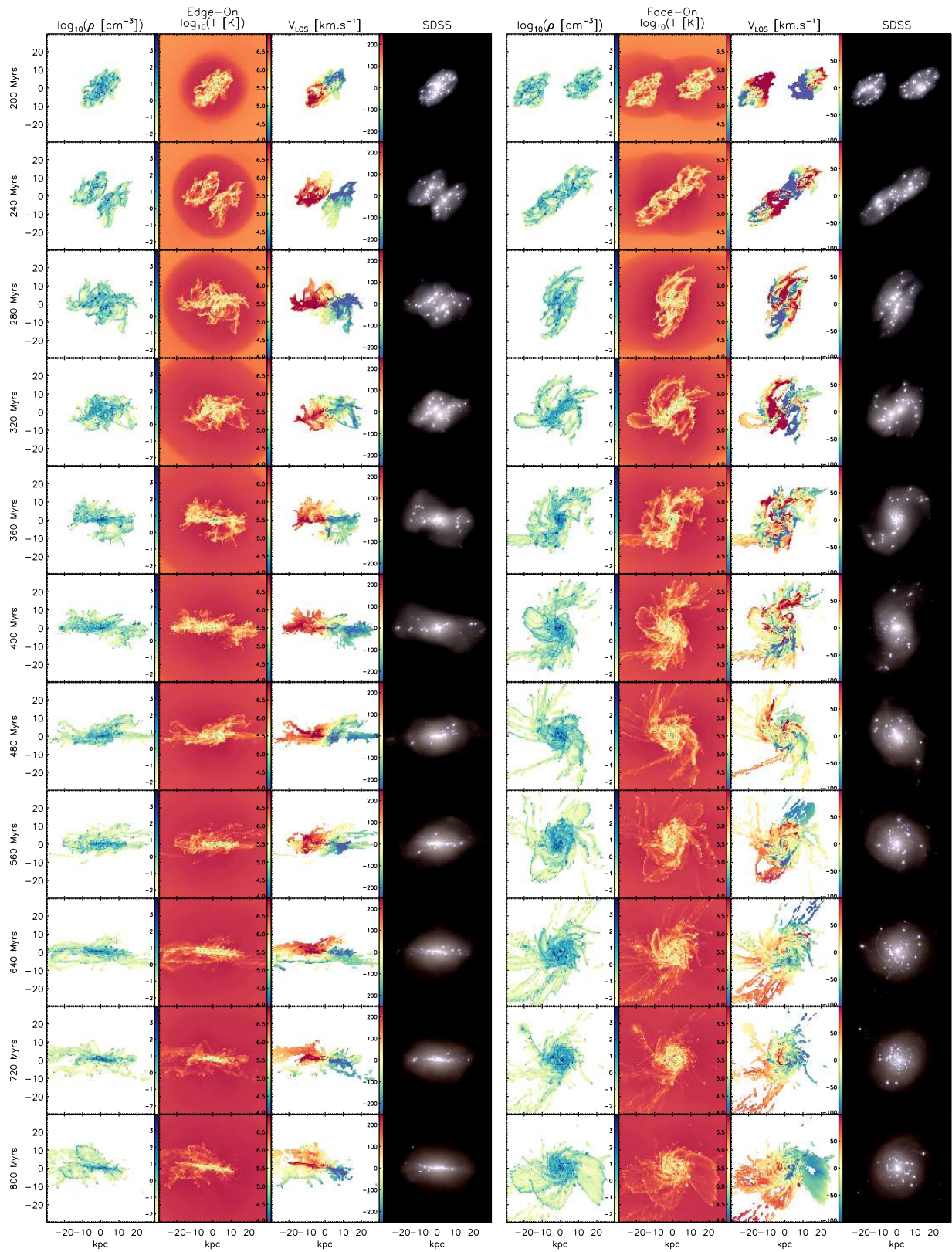
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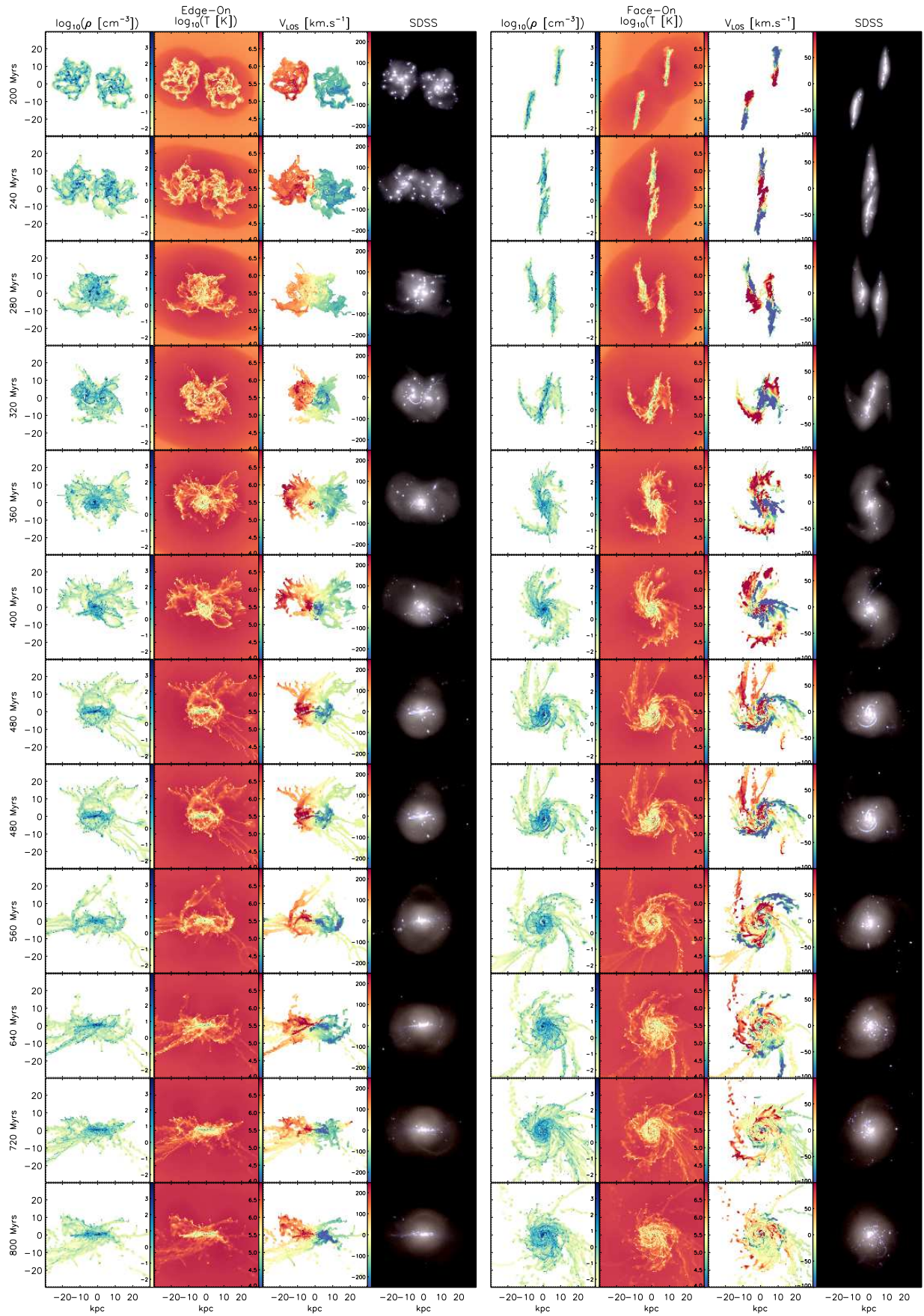
Suite A.2 : simulation G3.



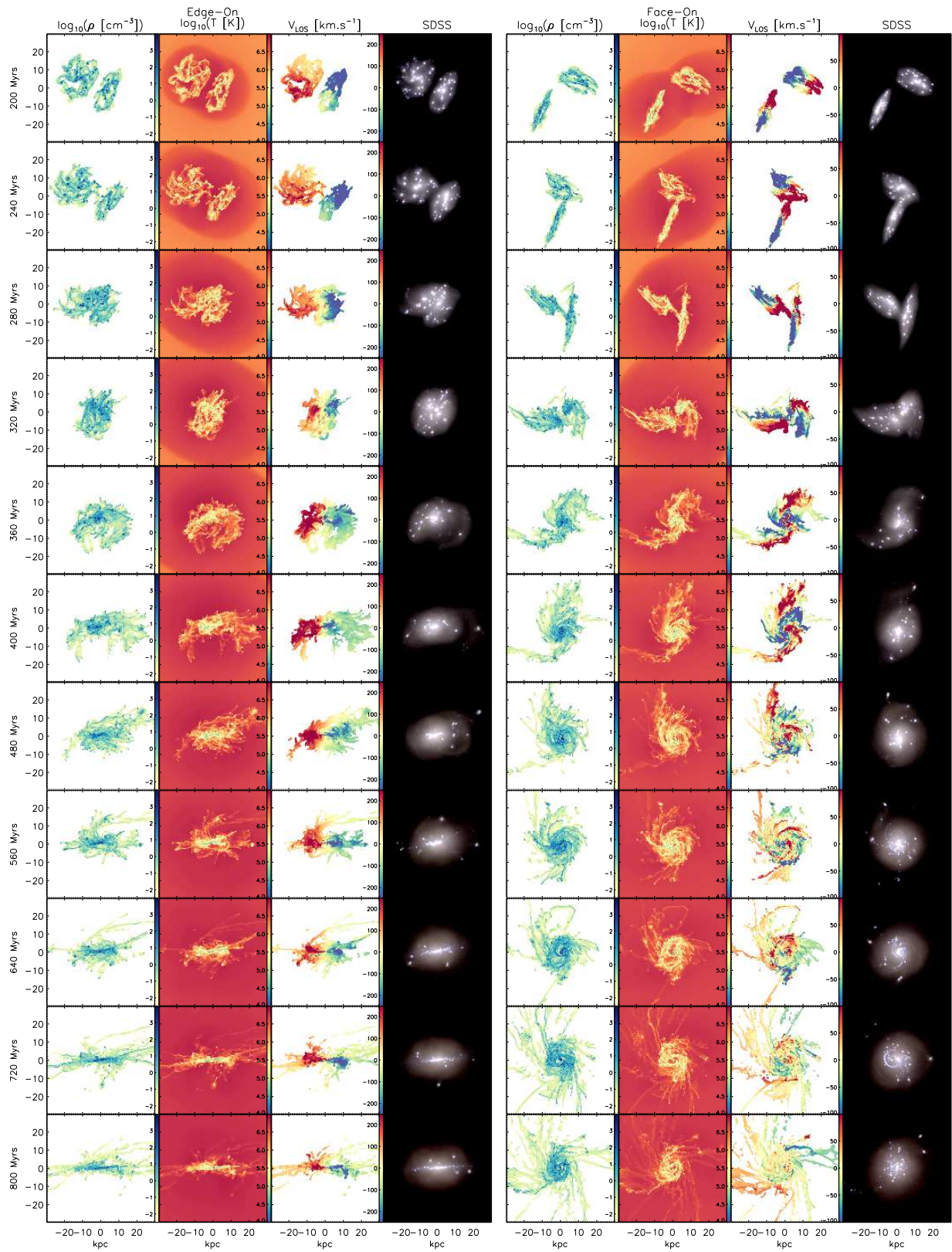
Suite A.2 : simulation G1_G1_90_0_90.



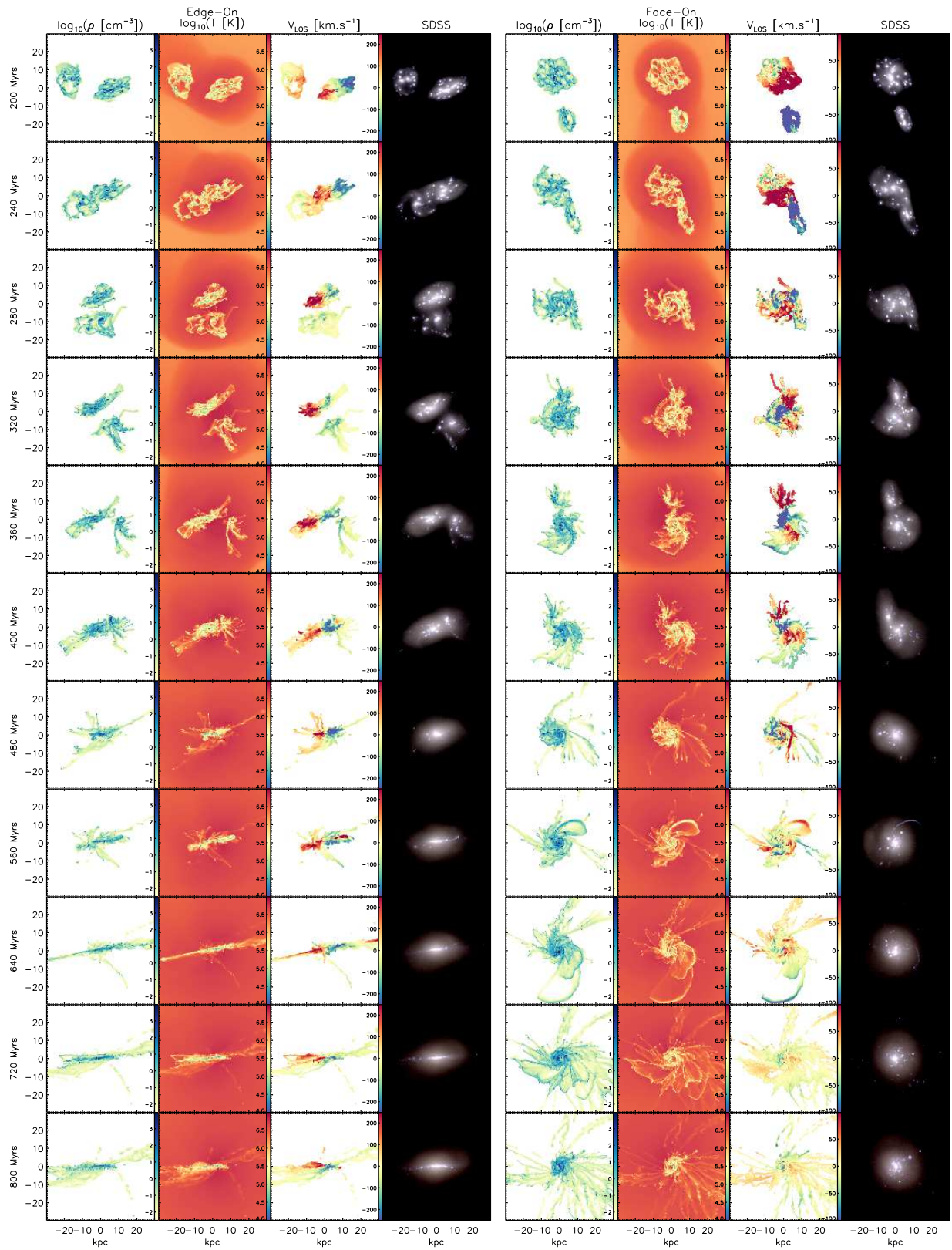
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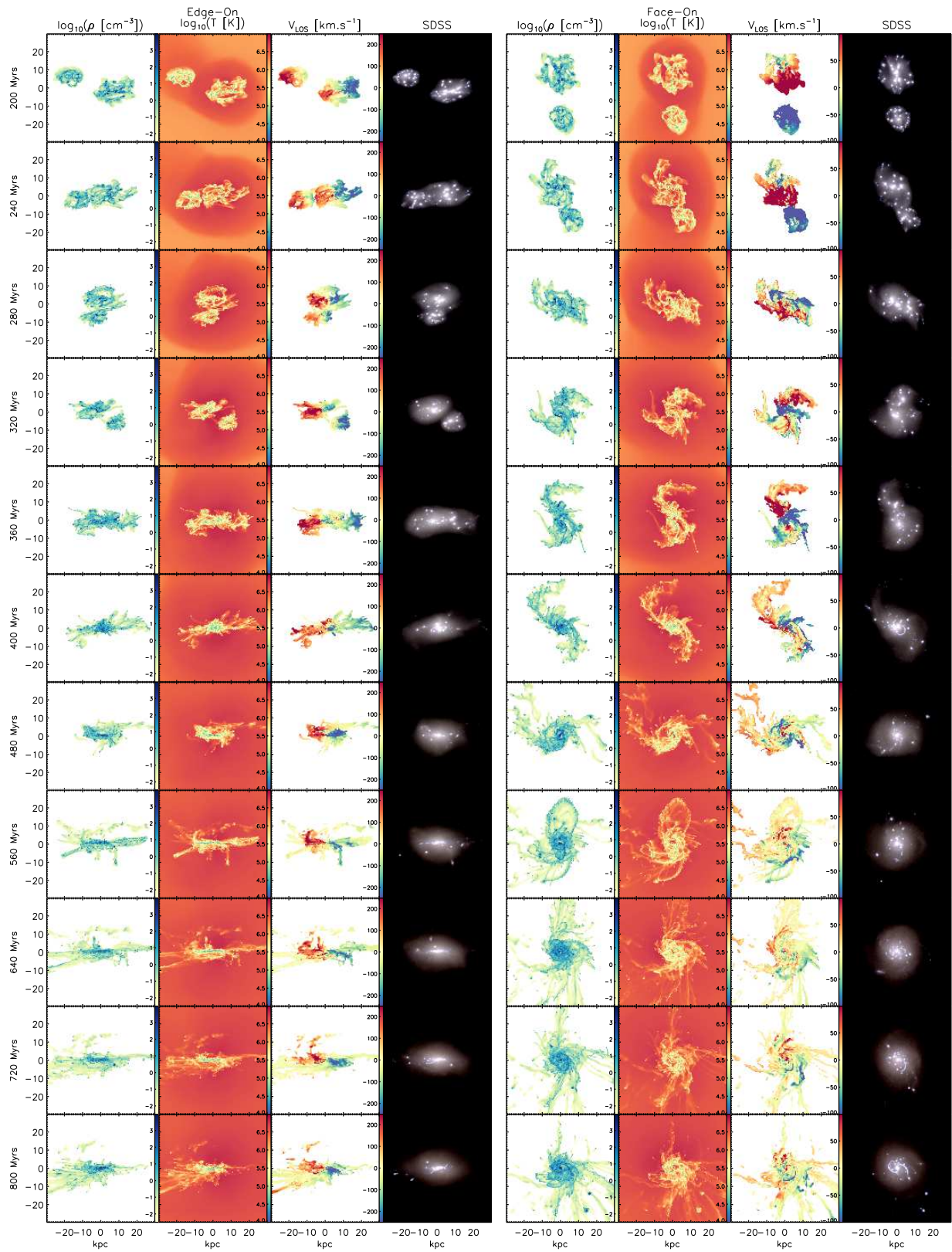
Suite A.2 : simulation G1_G1_90_90_180.



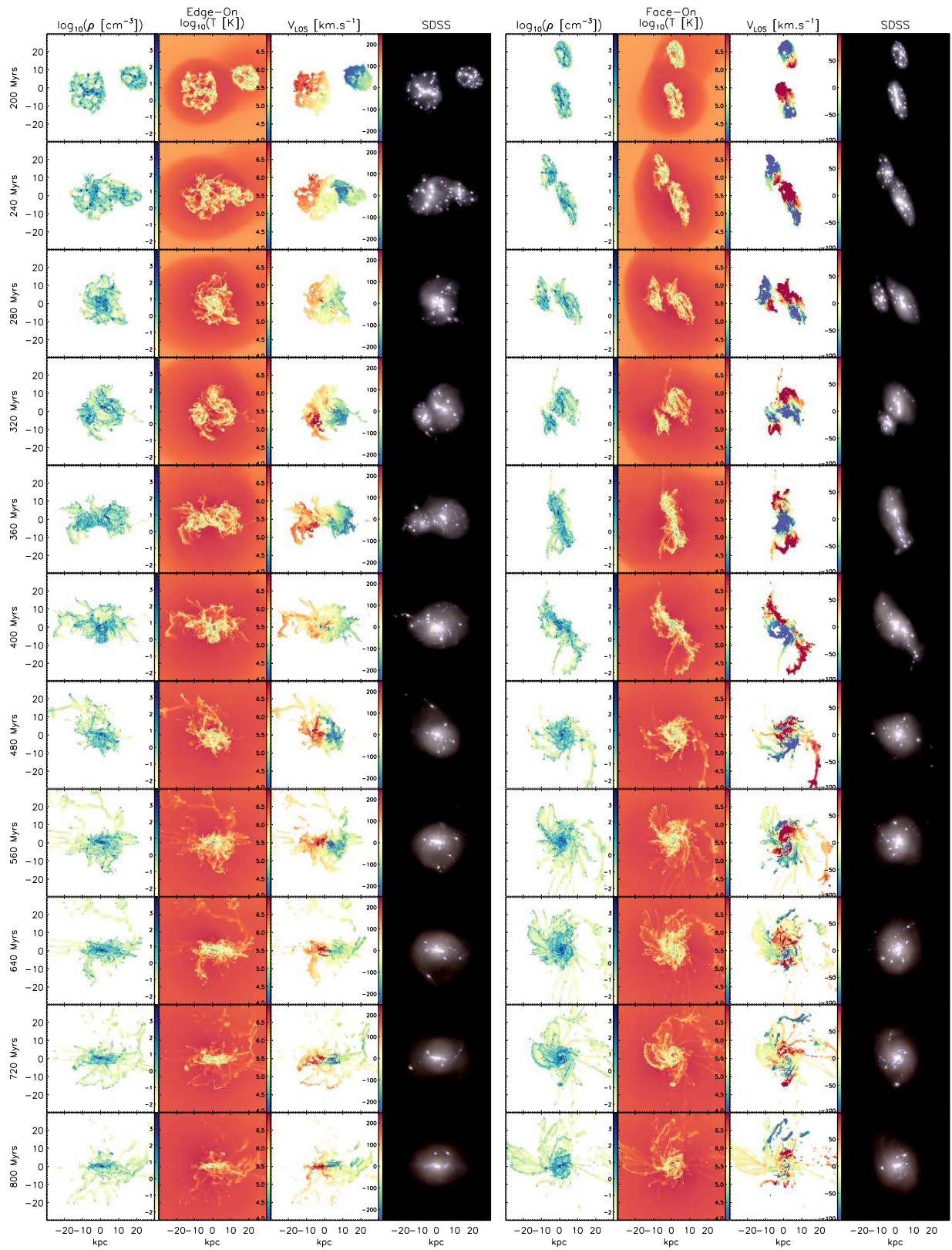
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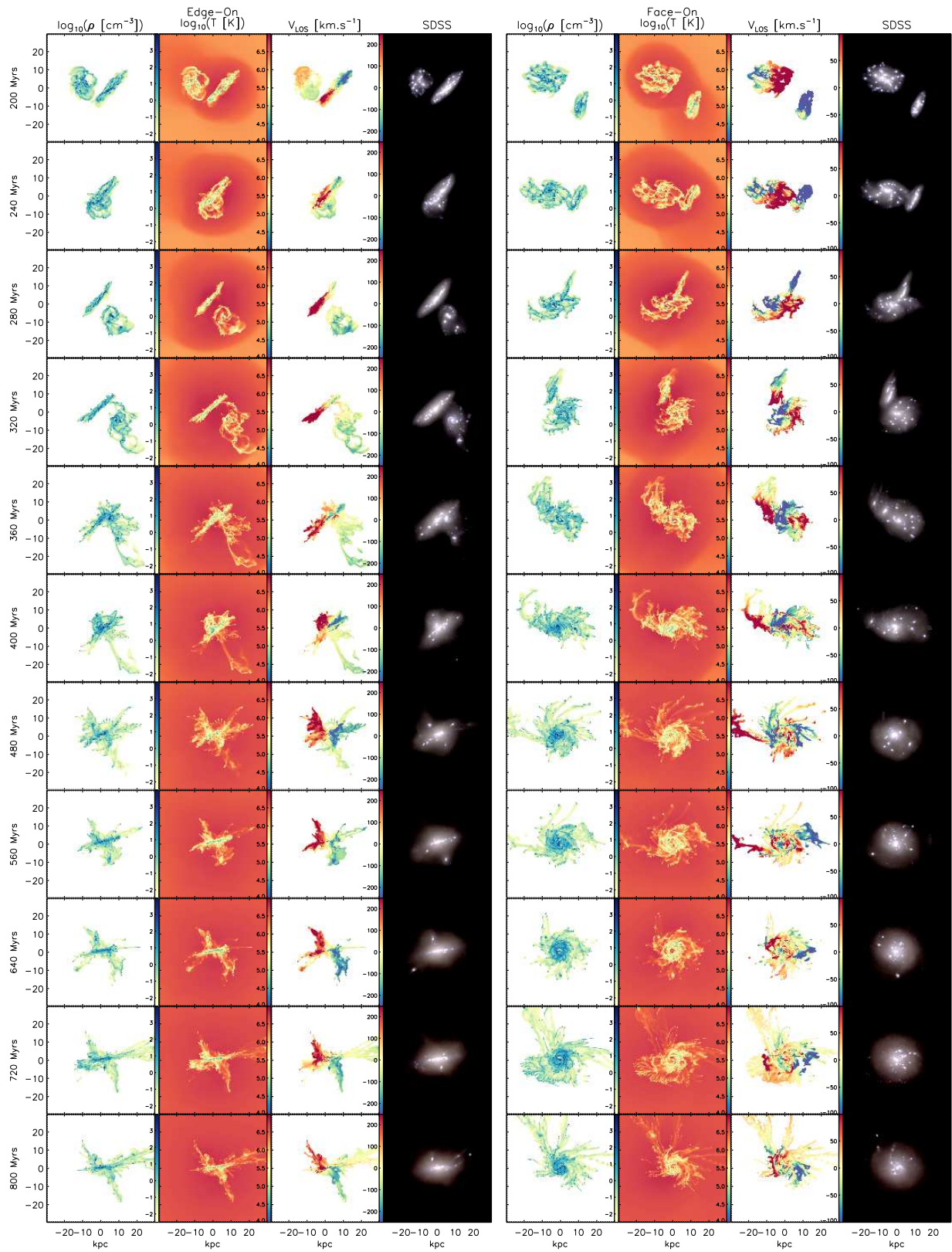
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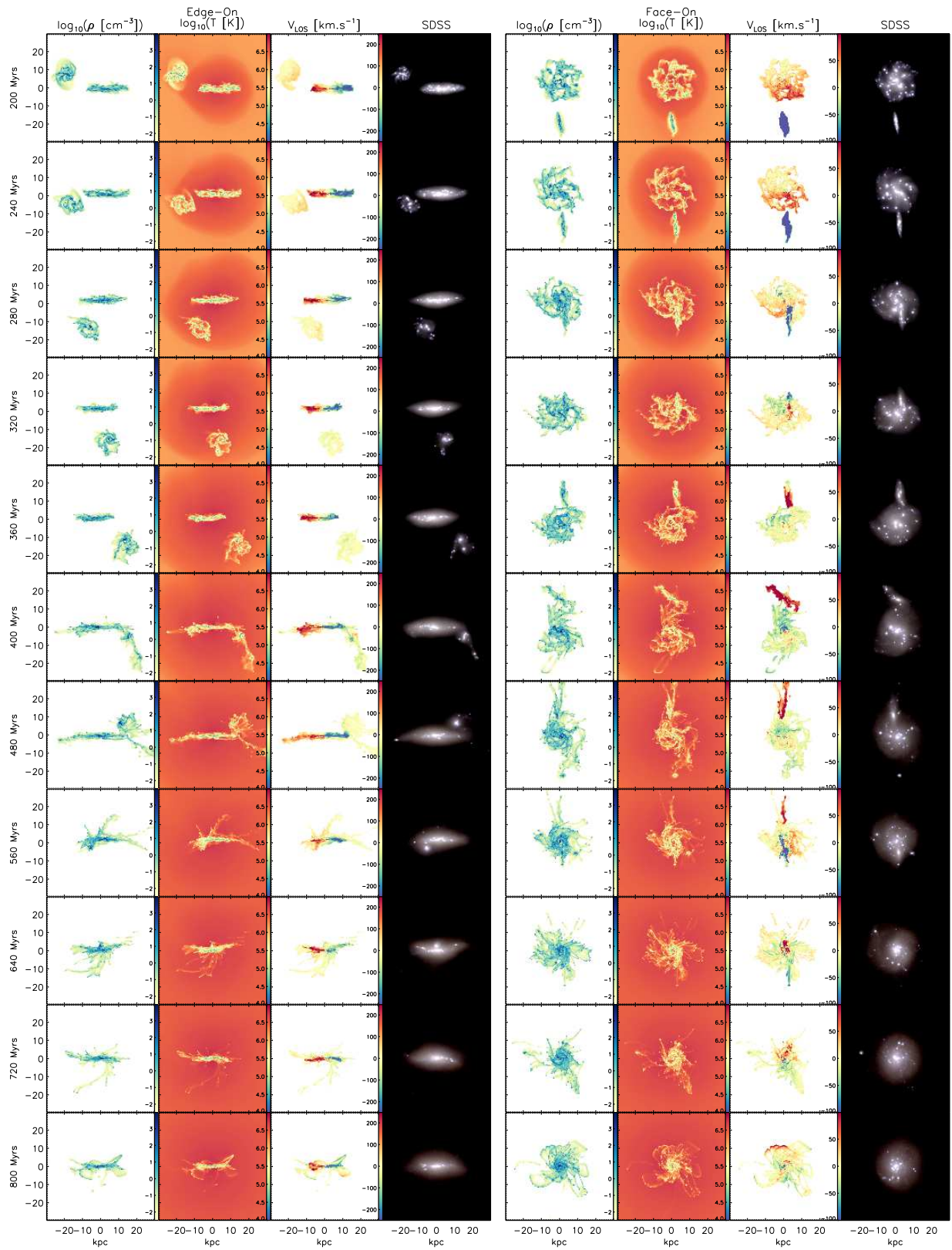
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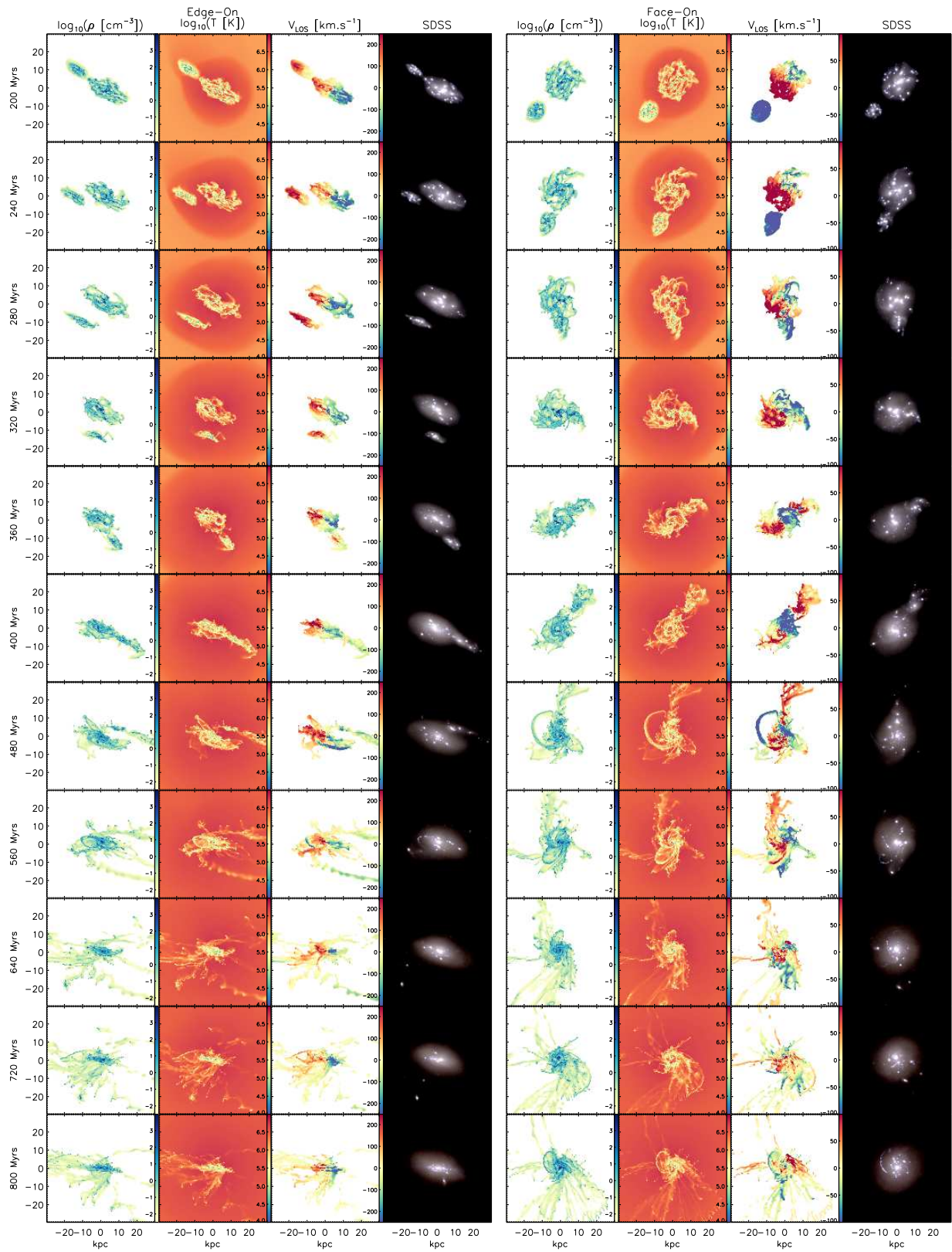
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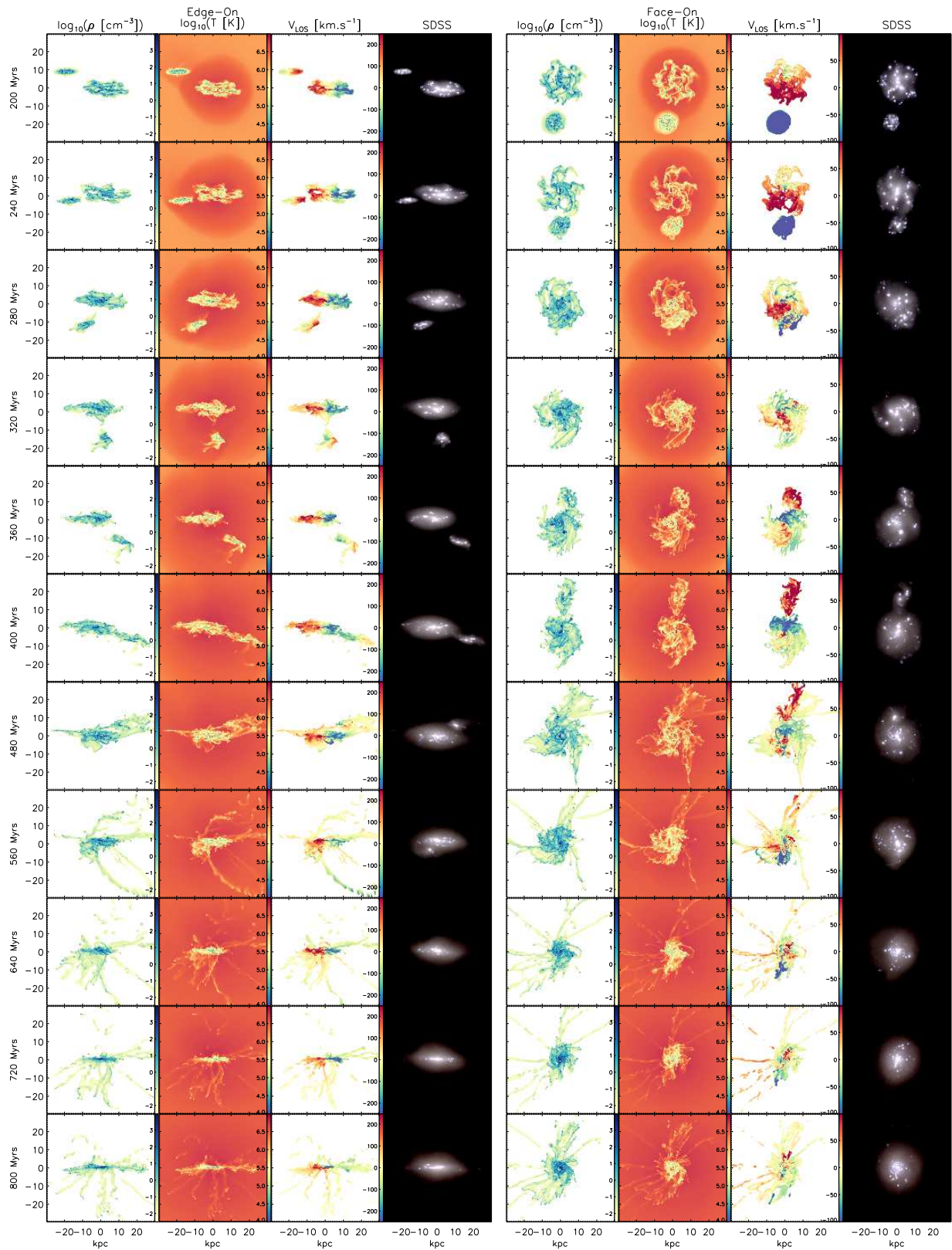
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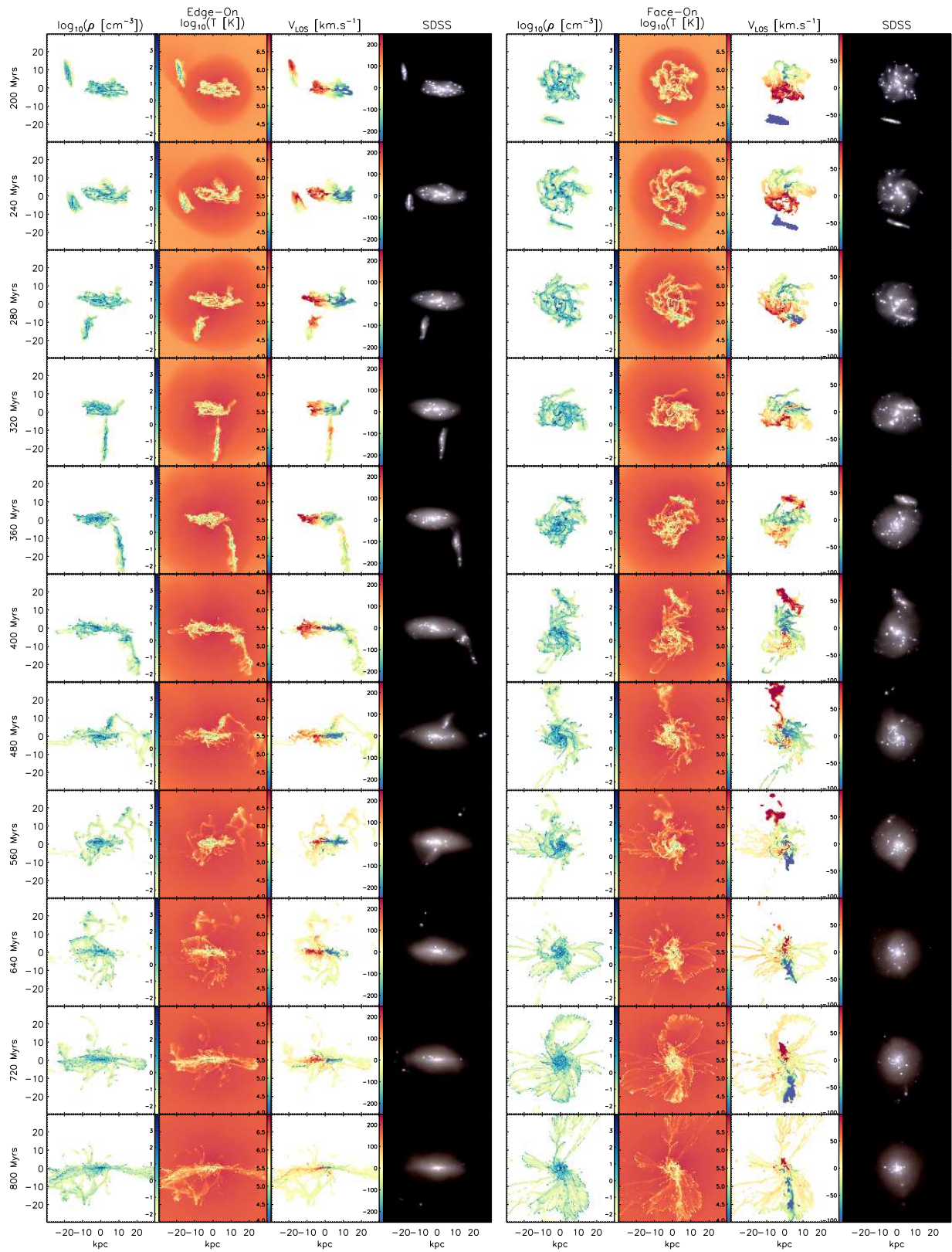
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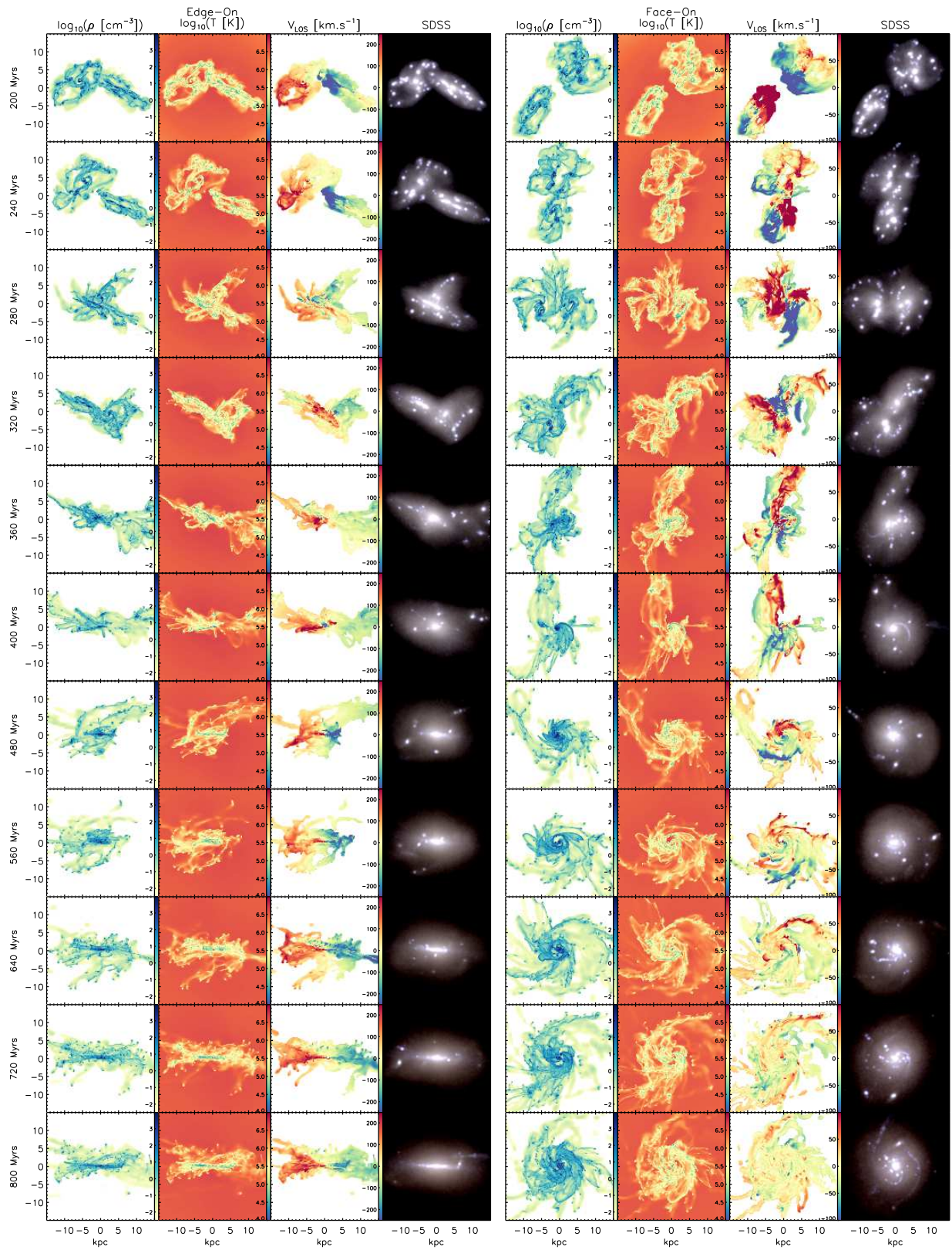
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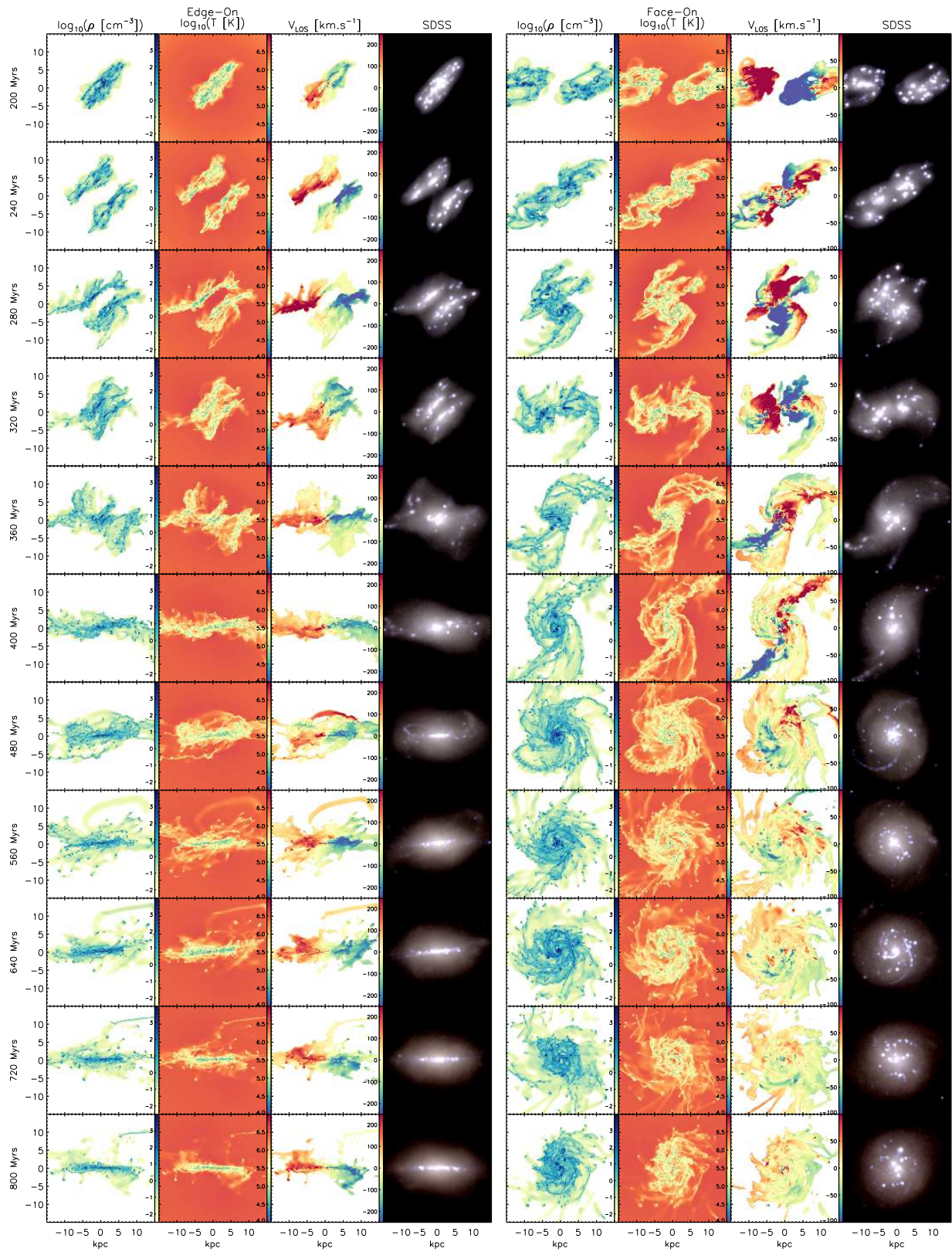
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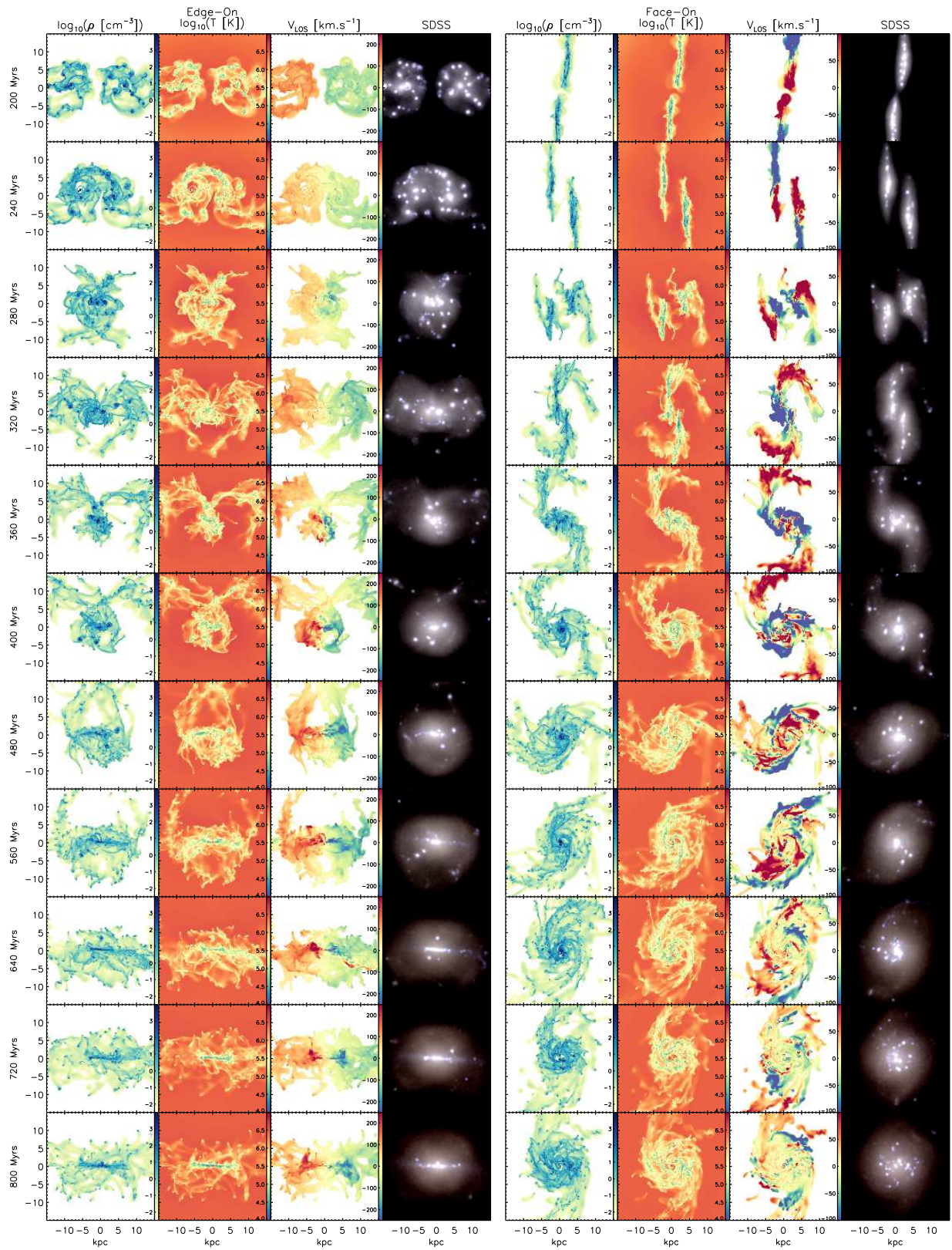
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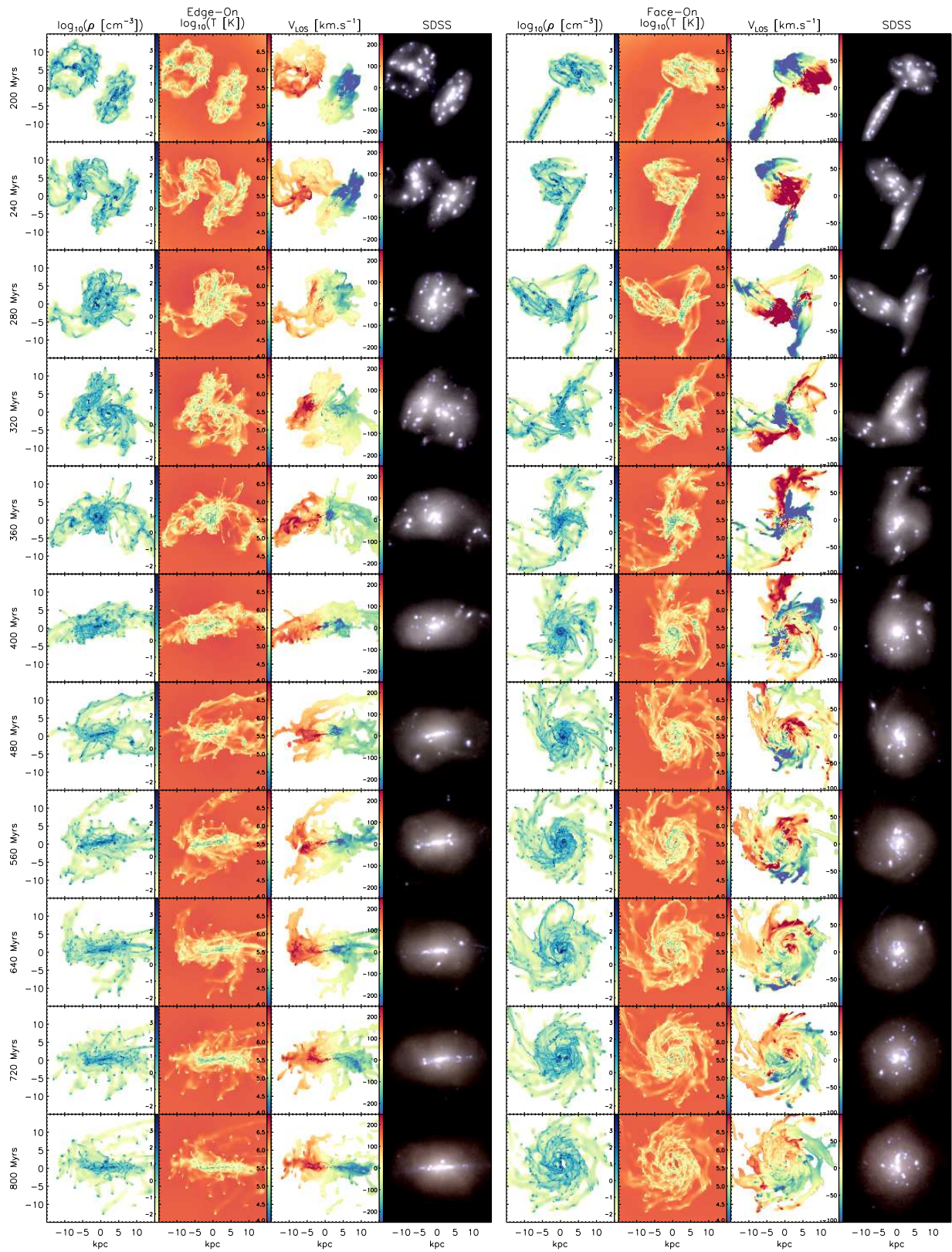
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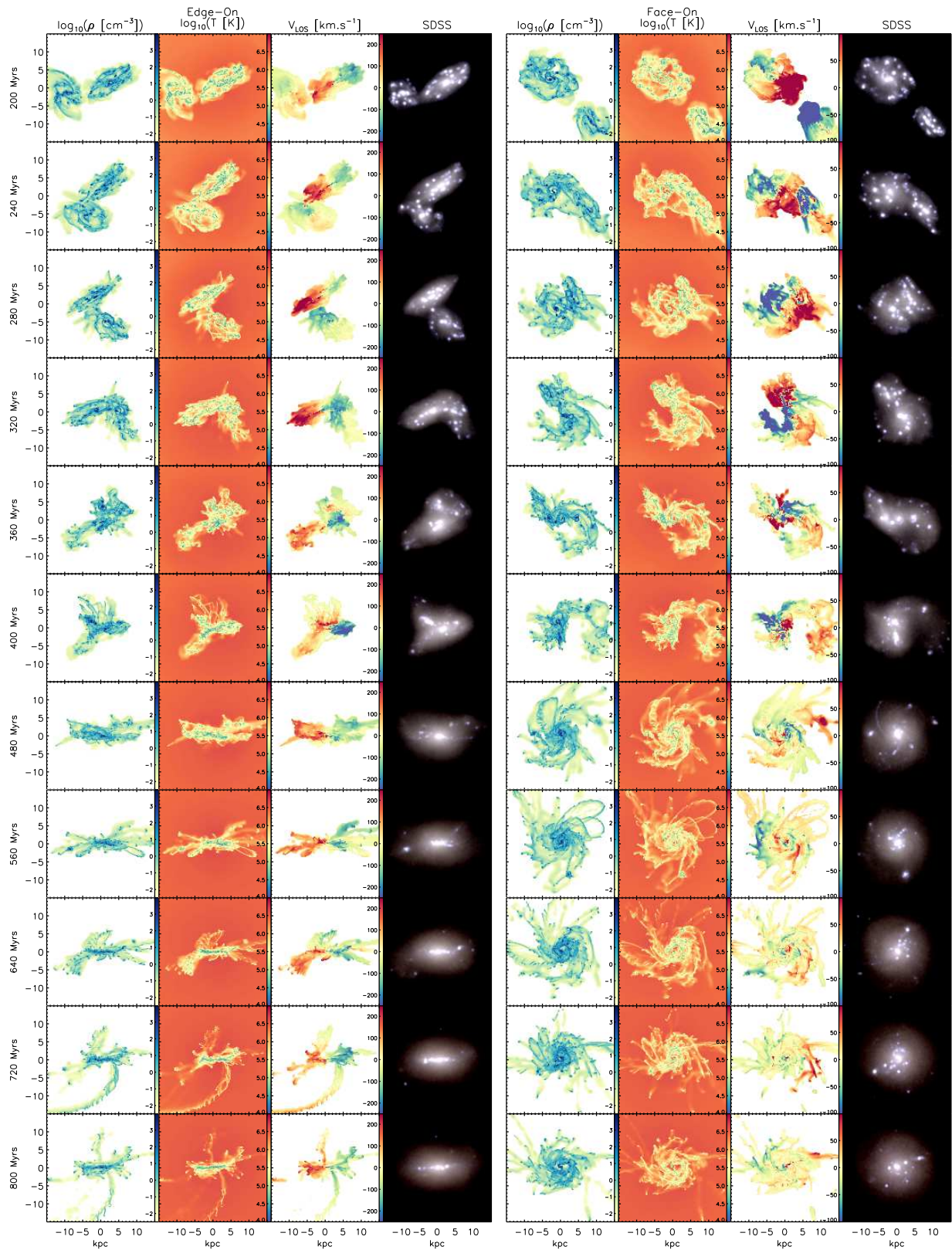
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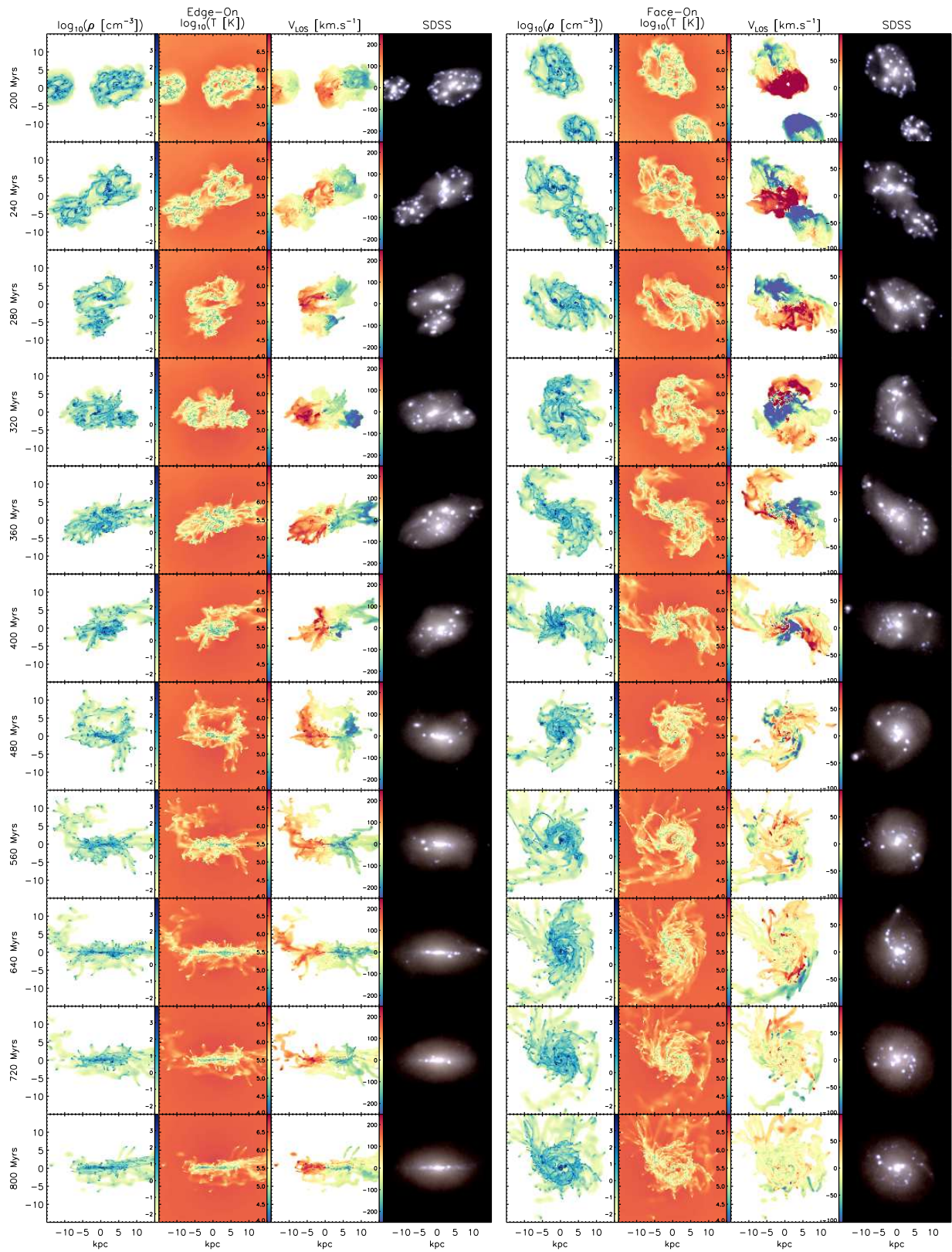
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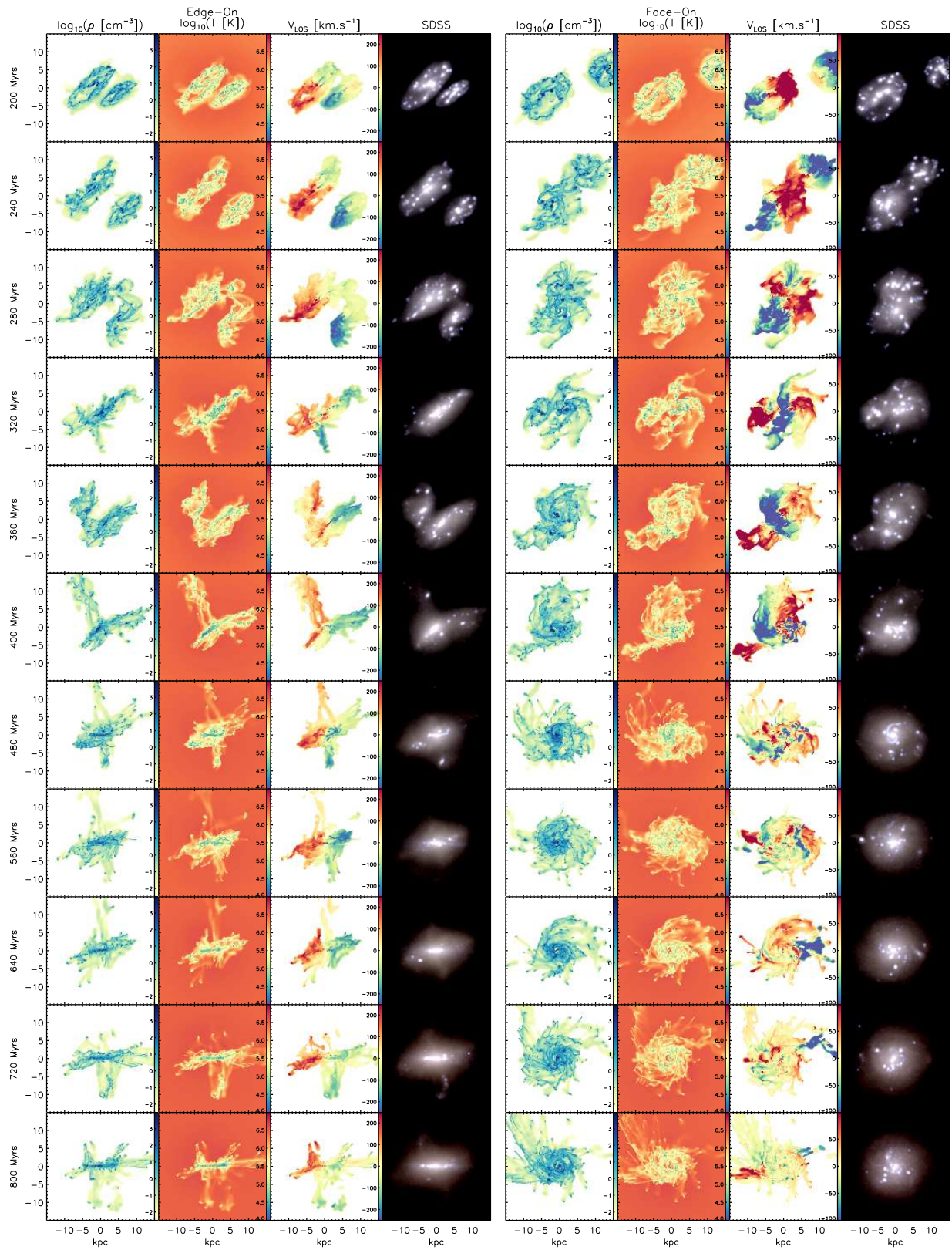
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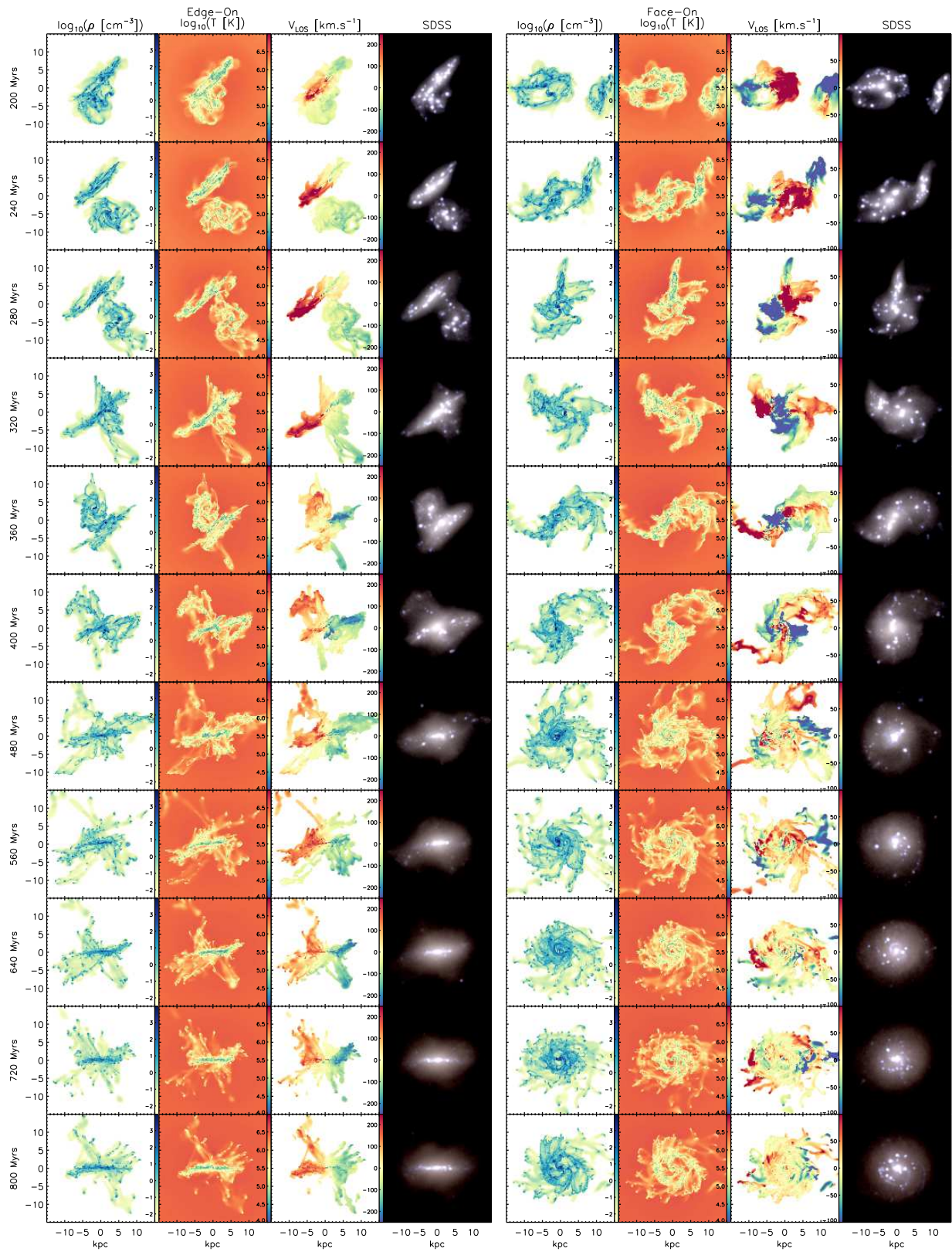
Suite A.2 : simulation G2_G3_90_0_90.



Suite A.2 : simulation G2_G3_90_90_0.



Suite A.2 : simulation G2_G3_90_90_180.



Suite A.2 : simulation G2_G3_90_90_90.

A.3 Champs de vitesses de pseudo-observations SINFONI [Chapitre 5]

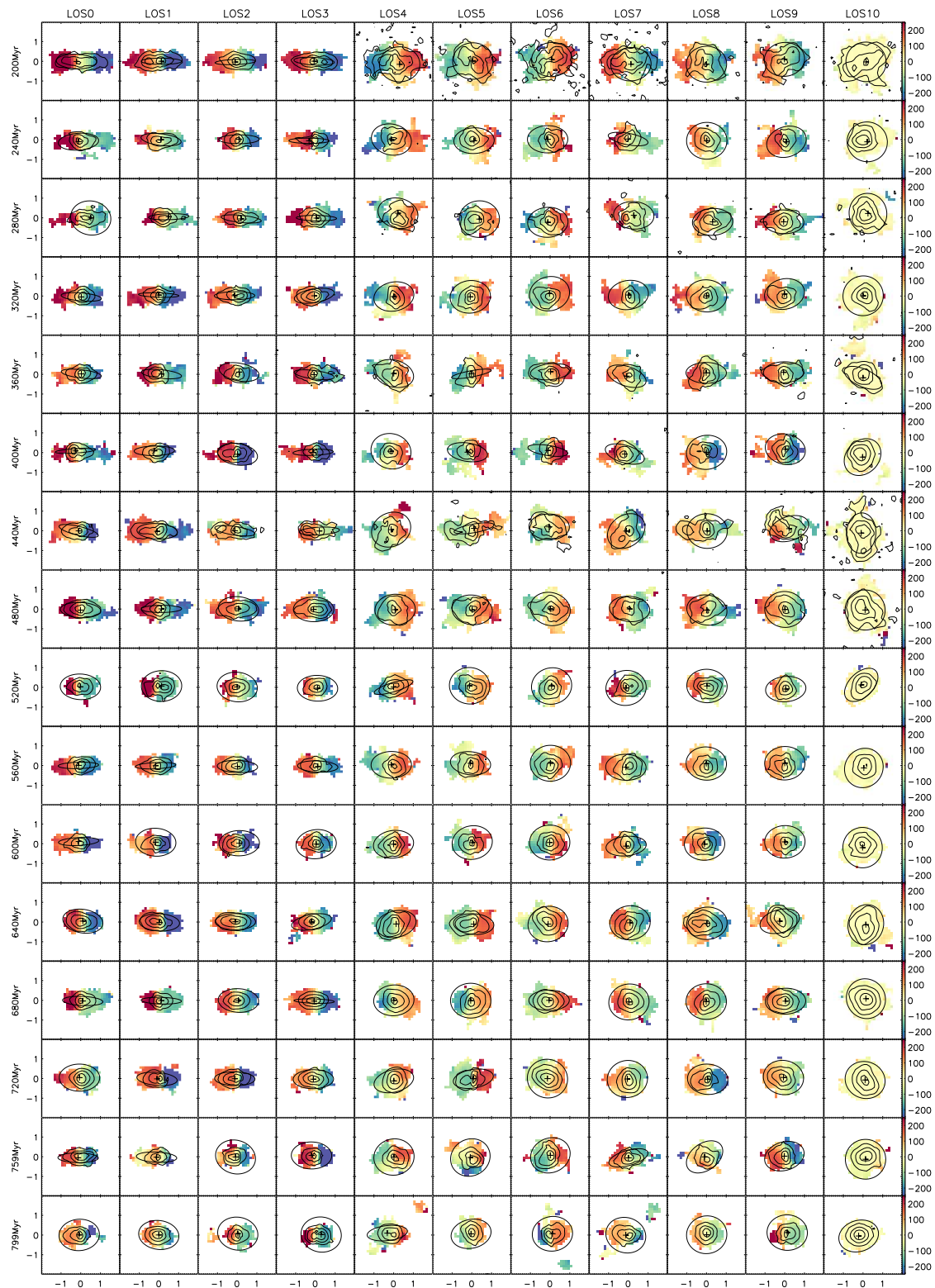
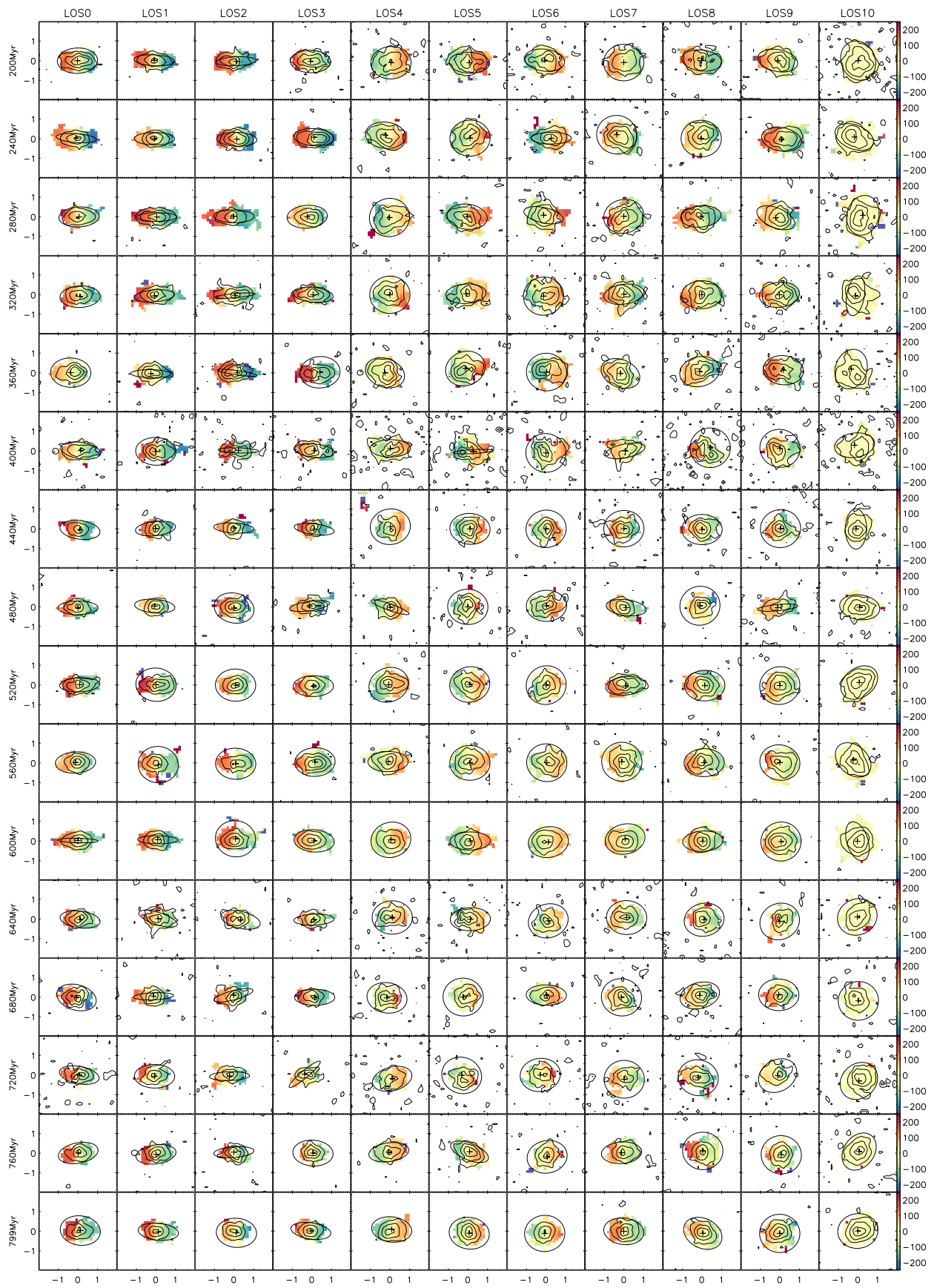
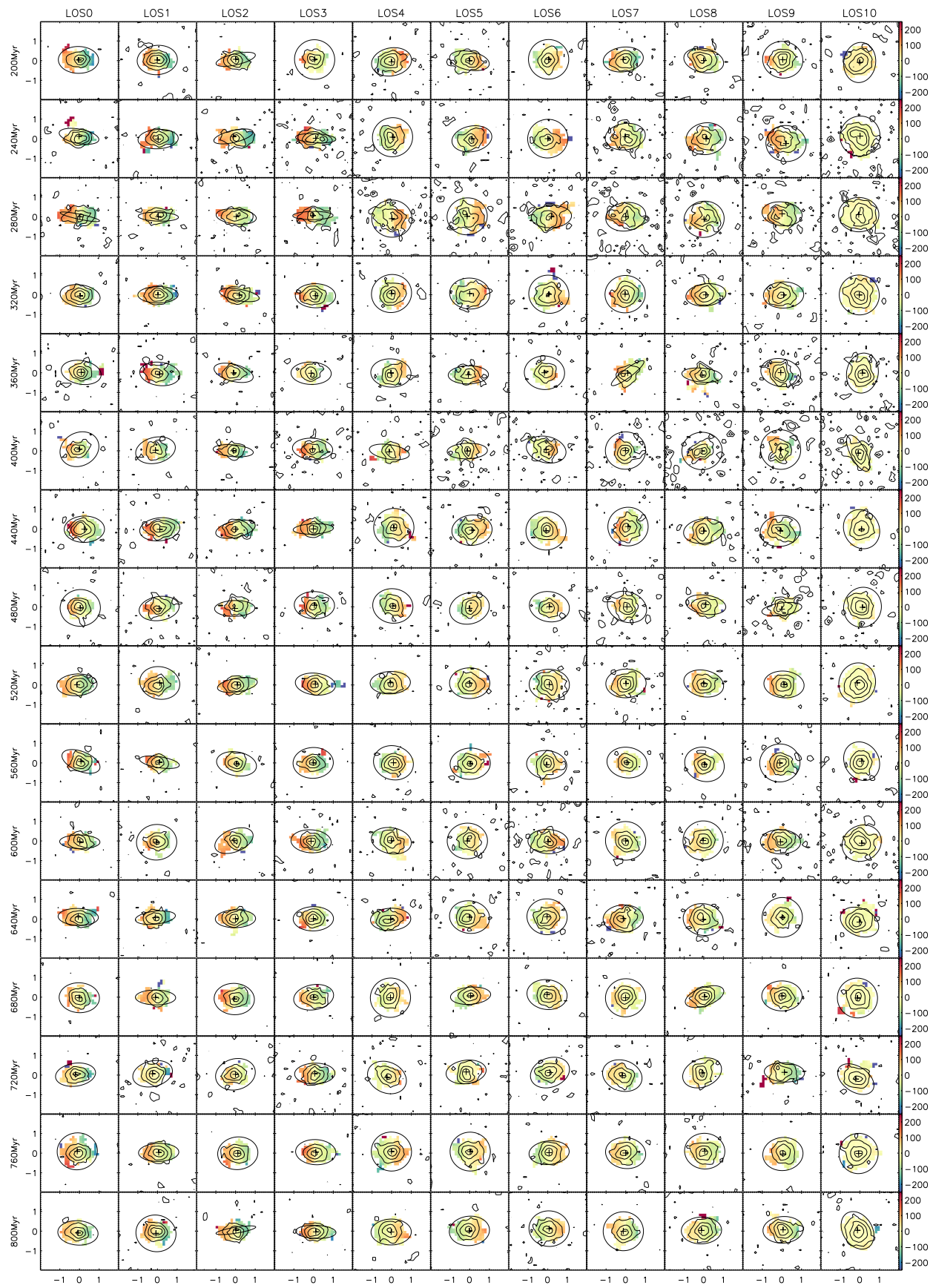


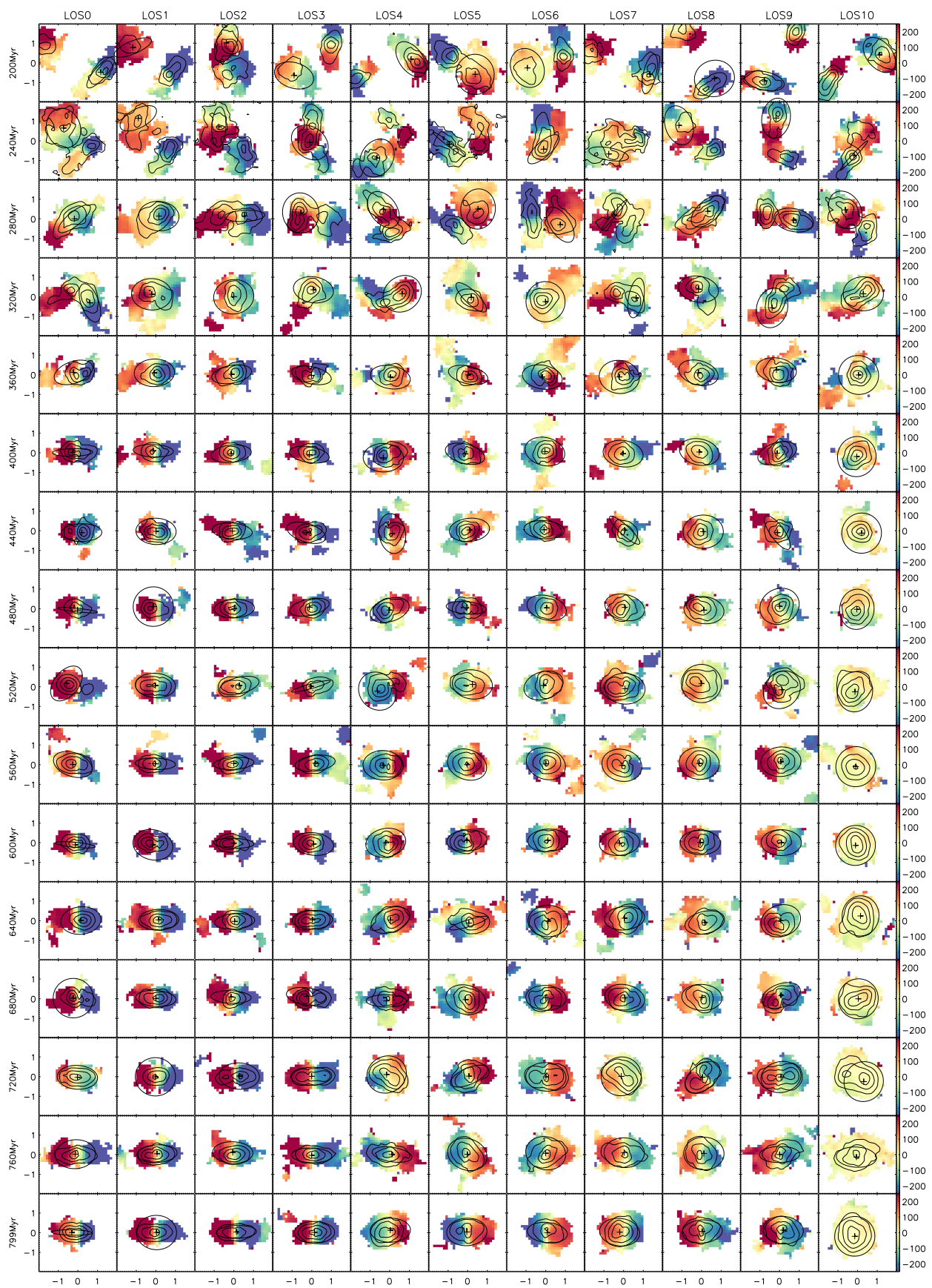
Figure A.3 – Champs de vitesse des pseudo-observations SINFONI de la simulation G1 de l'échantillon MIRAGE. Chaque ligne présente un temps spécifique dans la simulation, chaque colonne présente une ligne de visée différente pour la pseudo-observation. Sur chaque champ de vitesse, nous traçons les contours issue de la carte de flux $H\alpha$. Il est aussi tracé une ellipse centrée sur une croix afin de présenter le résultat de l'ajustement des paramètres géométrique grâce à la routine KINEGEN.



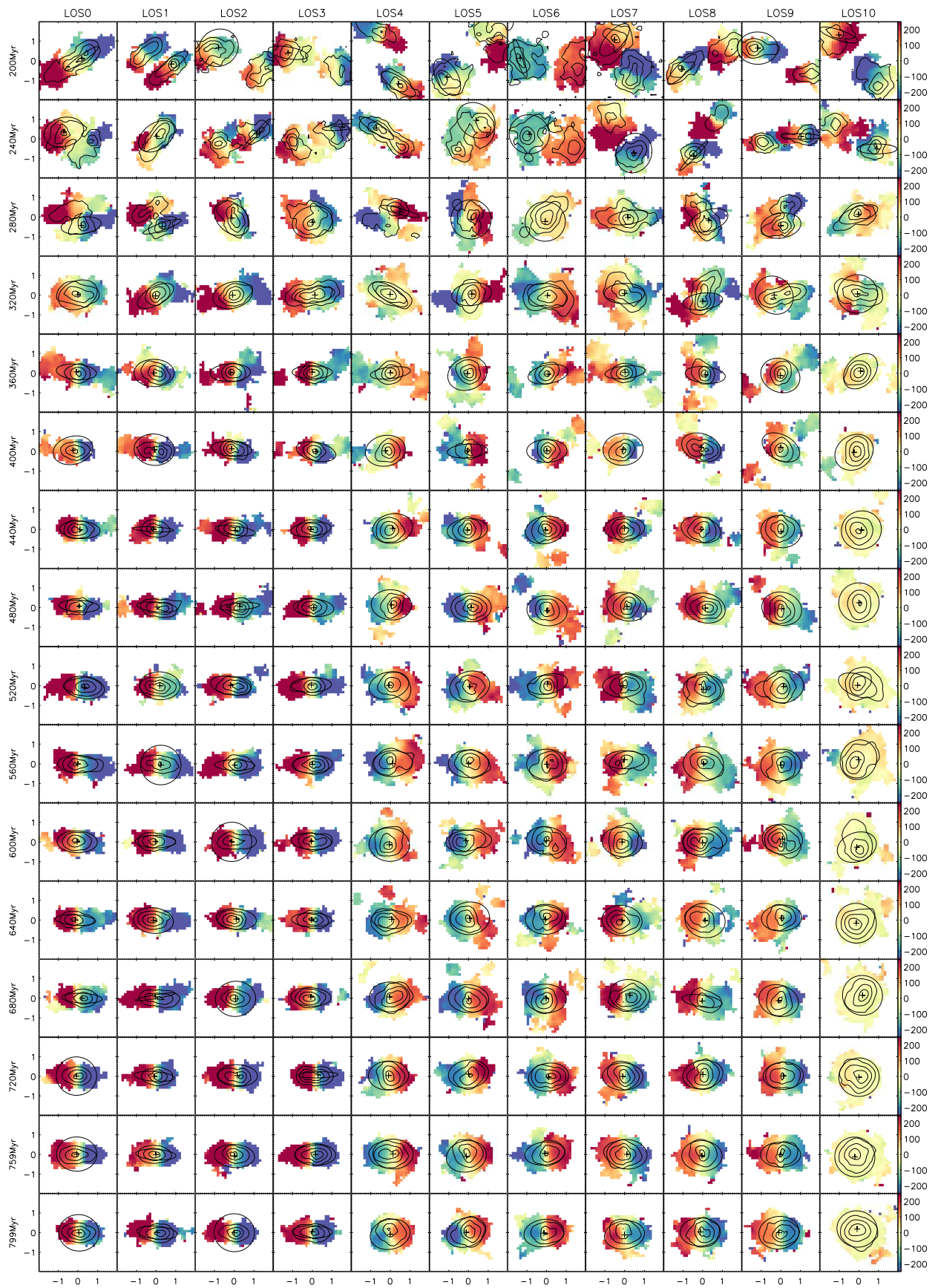
Suite A.3 : simulation de disque isolé G2.



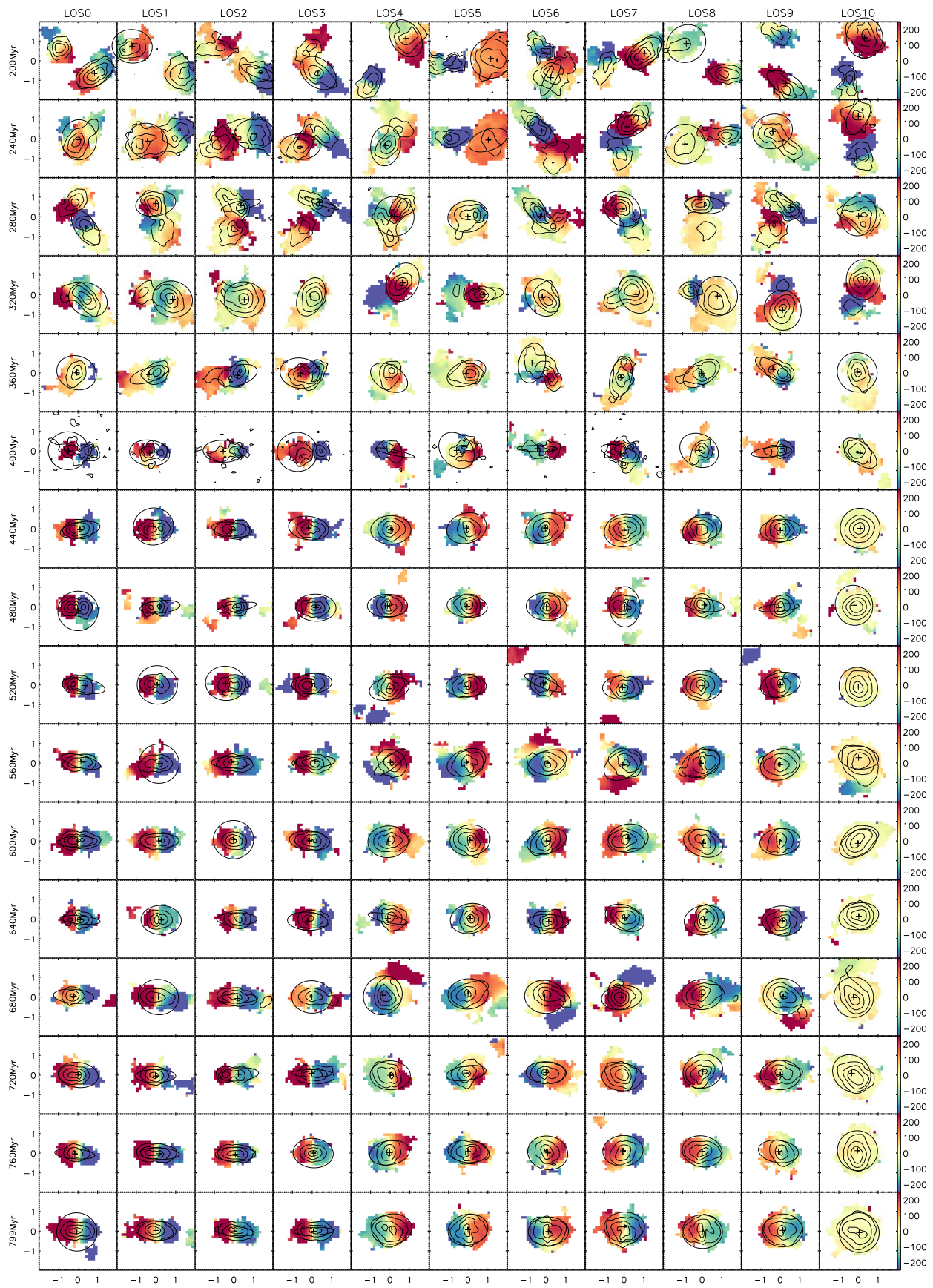
Suite A.3 : simulation de disque isolé G3.



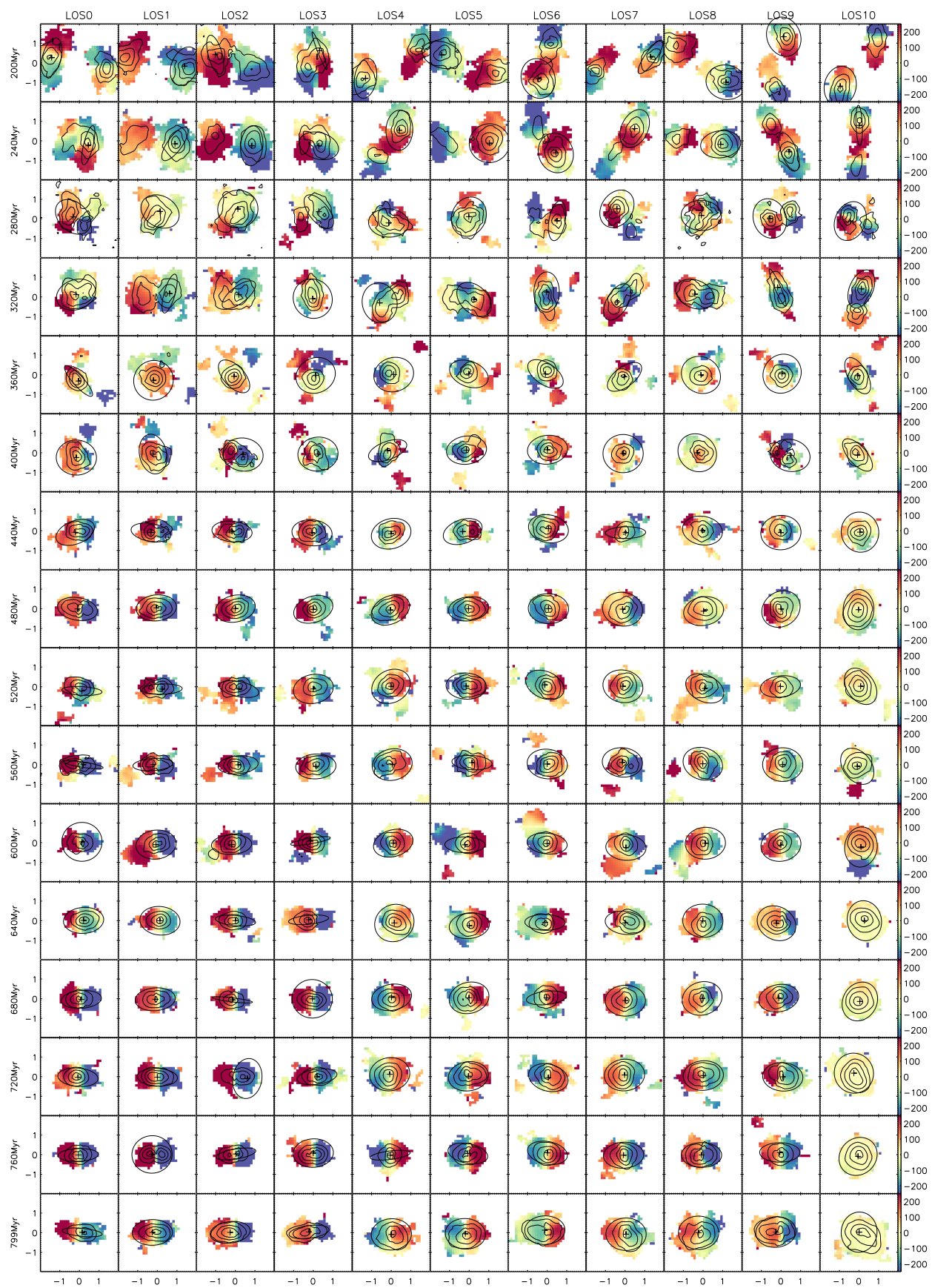
Suite A.3 : simulation de fusion G1_G1_90_90_90.



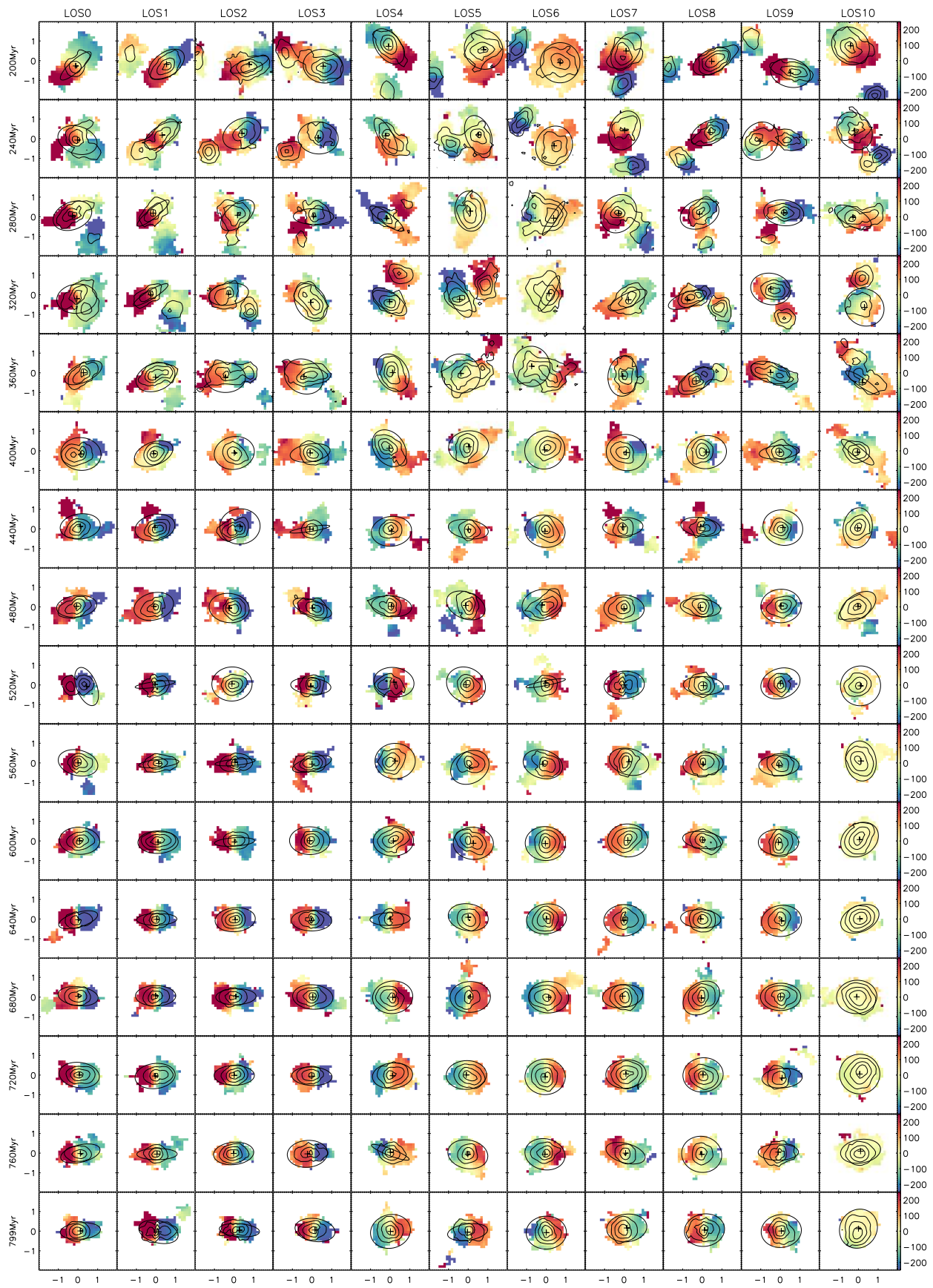
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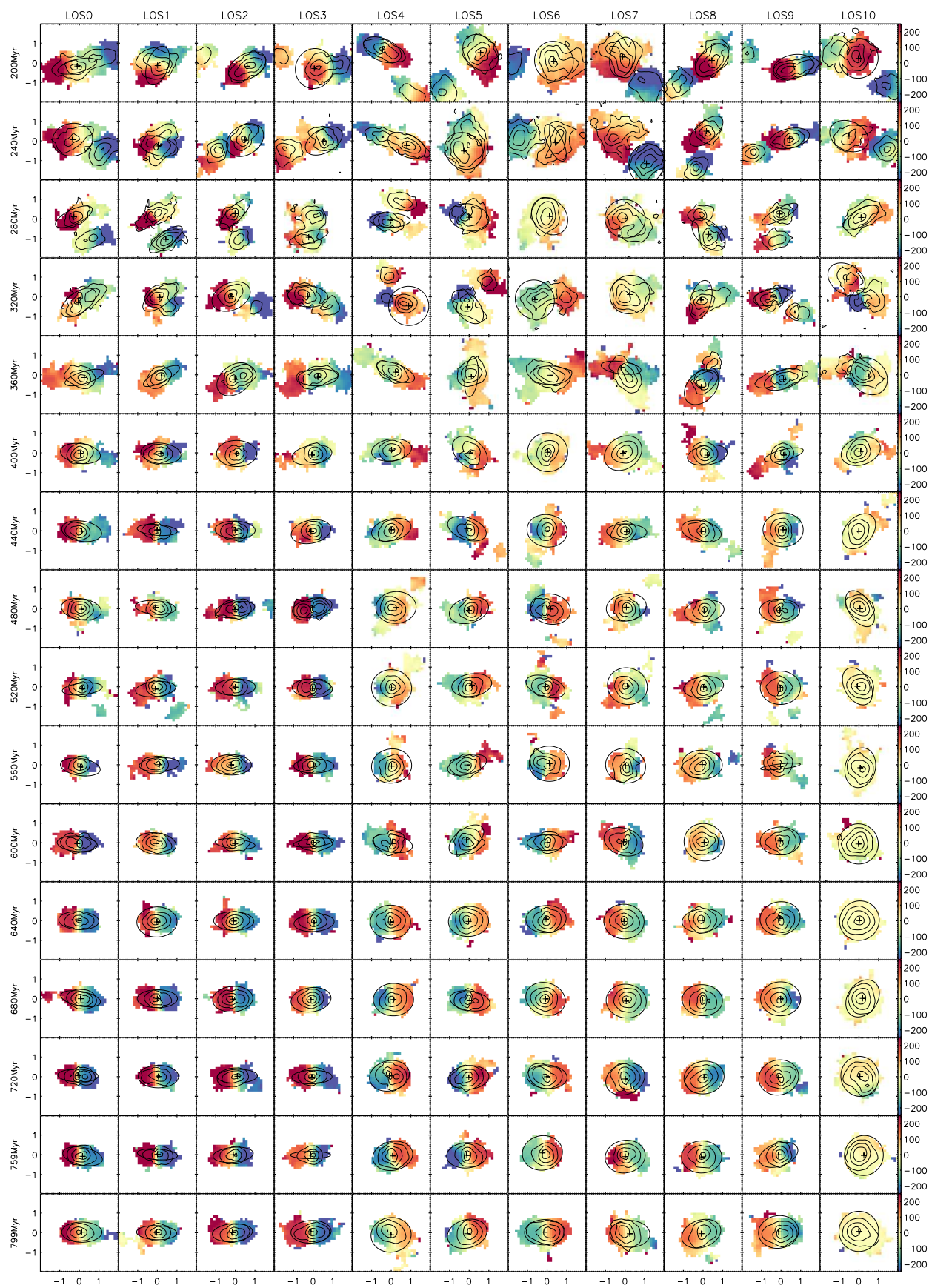
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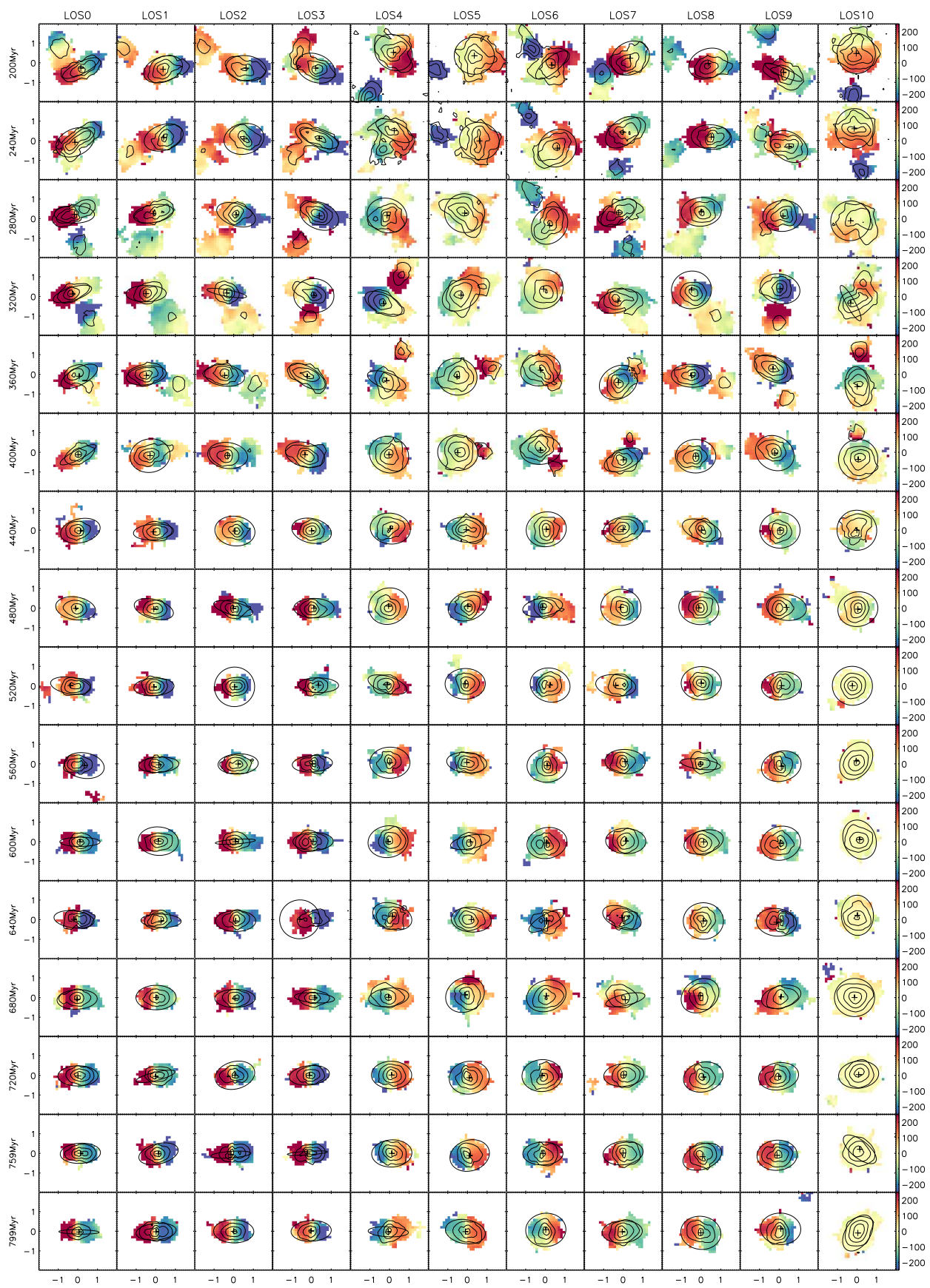
Suite A.3 : simulation de fusion G1_G1_90_90_180.



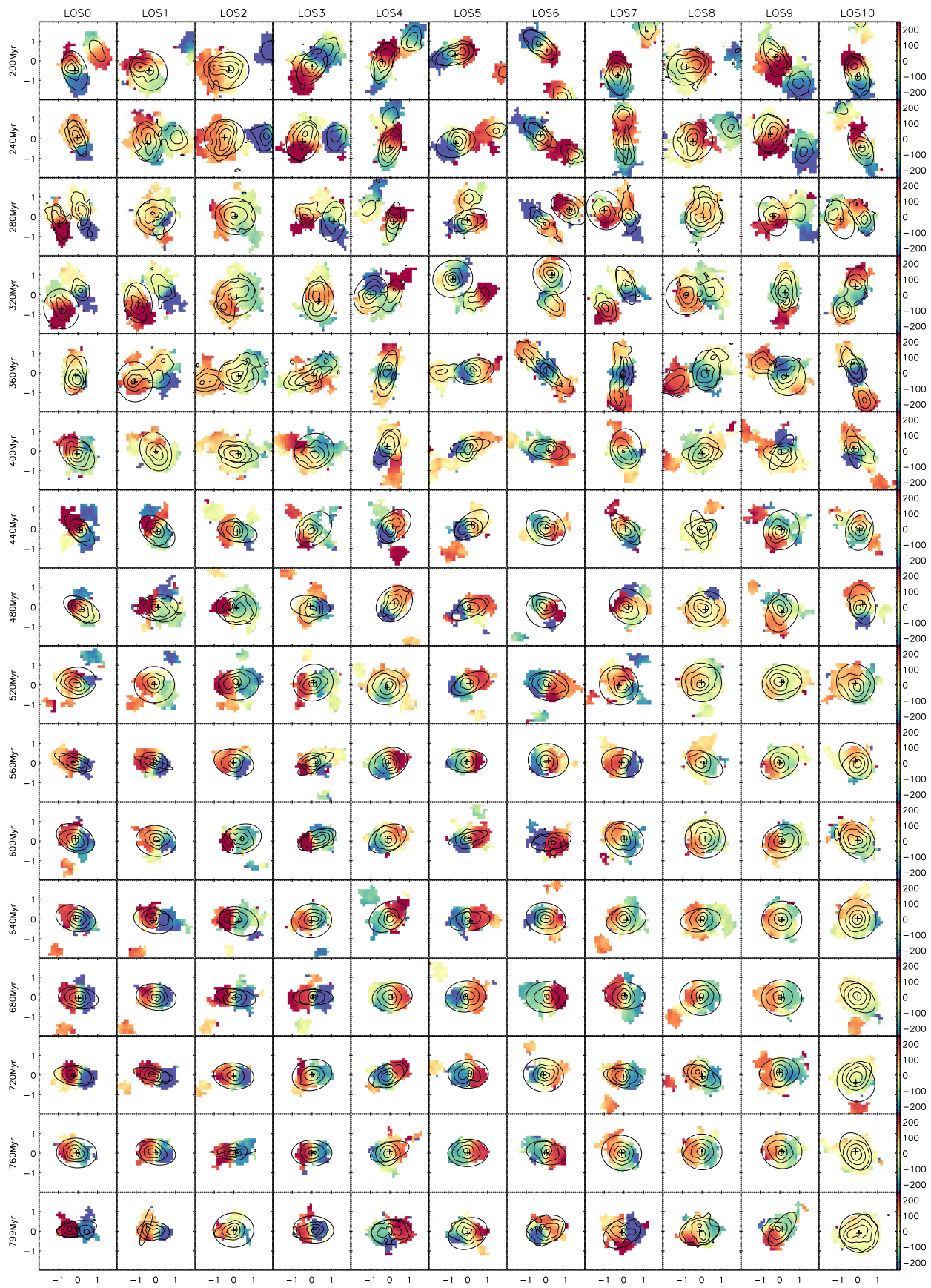
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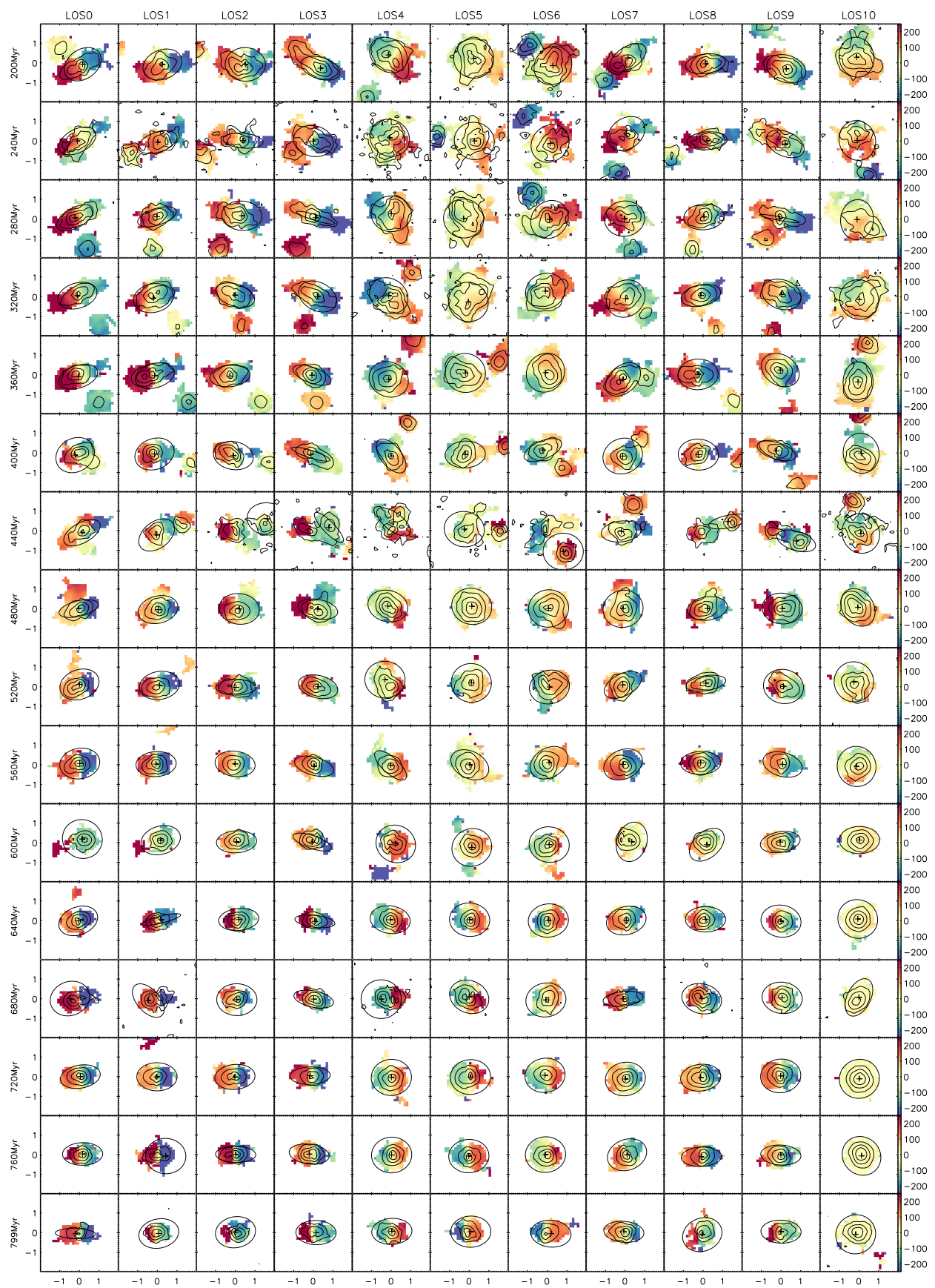
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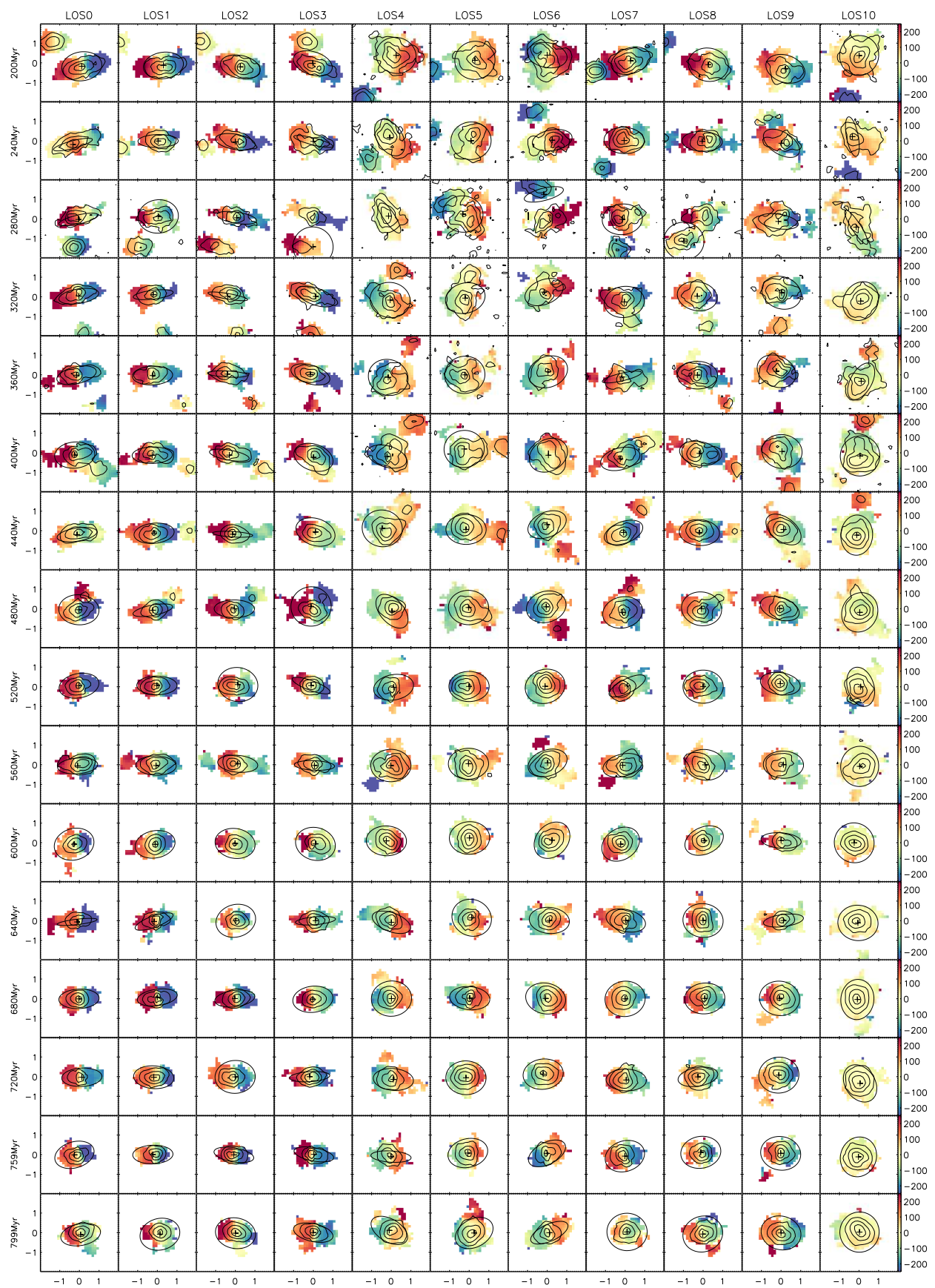
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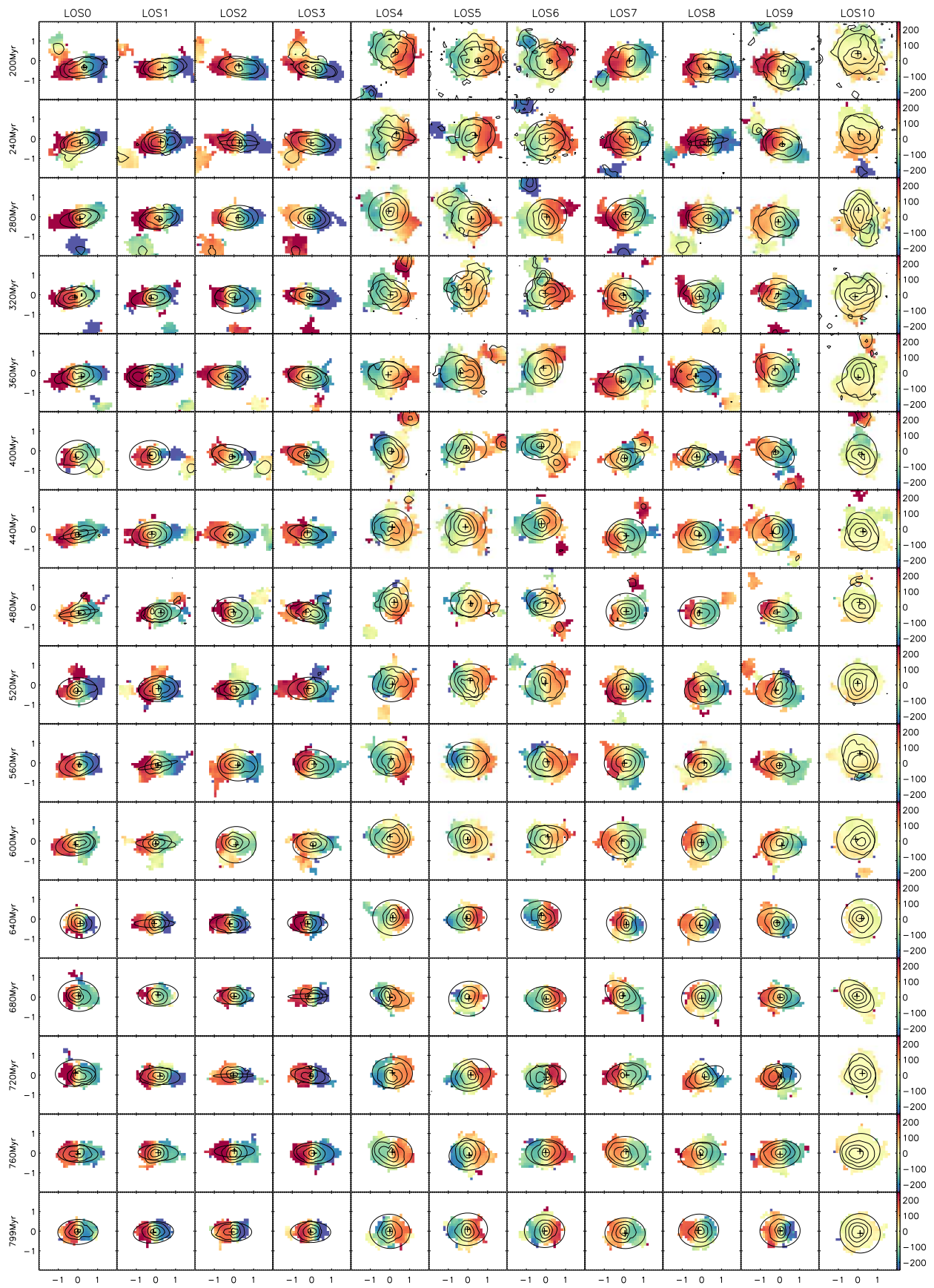
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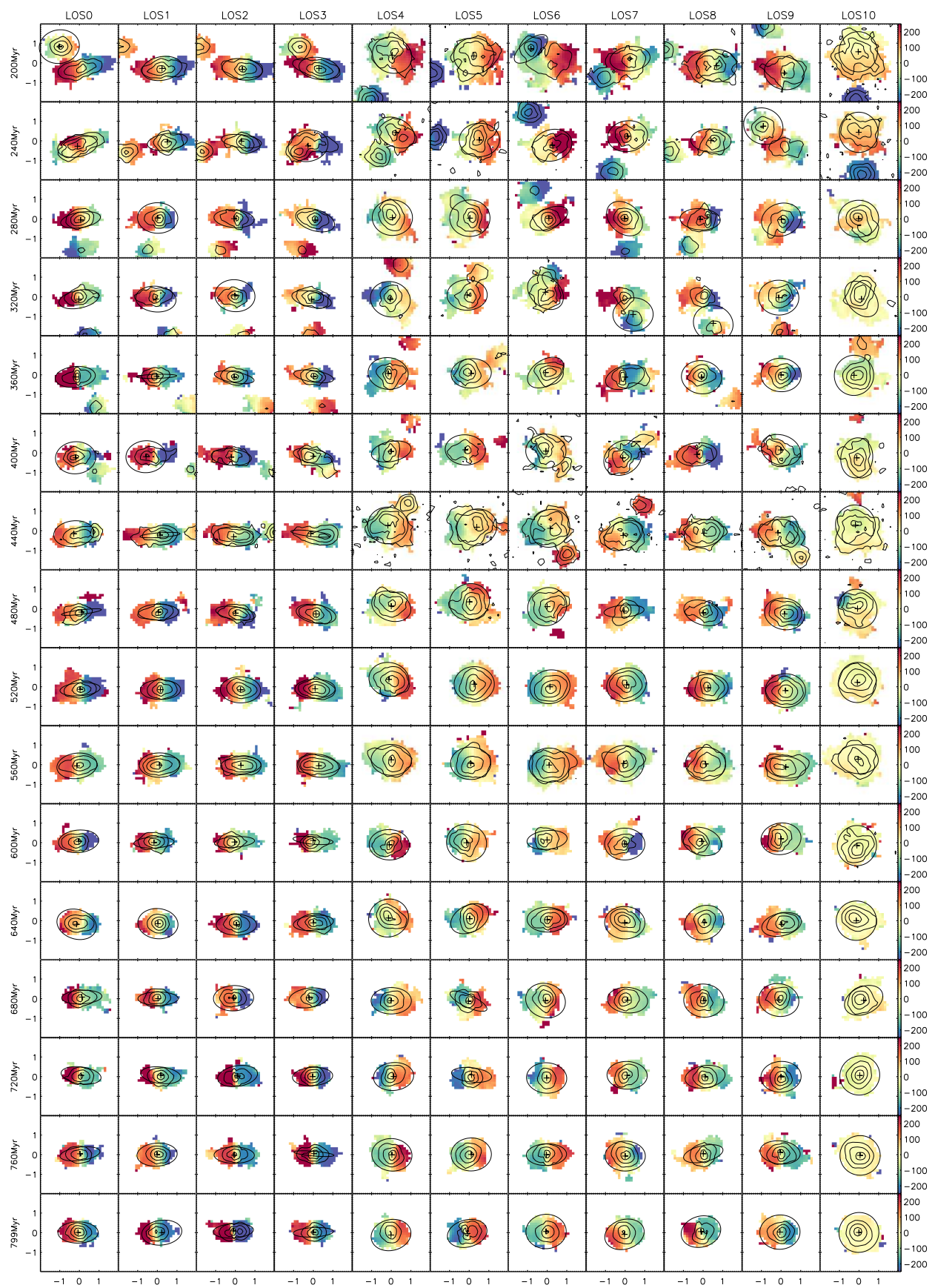
Suite A.3 : simulation de fusion G1_G3_90_90_90.



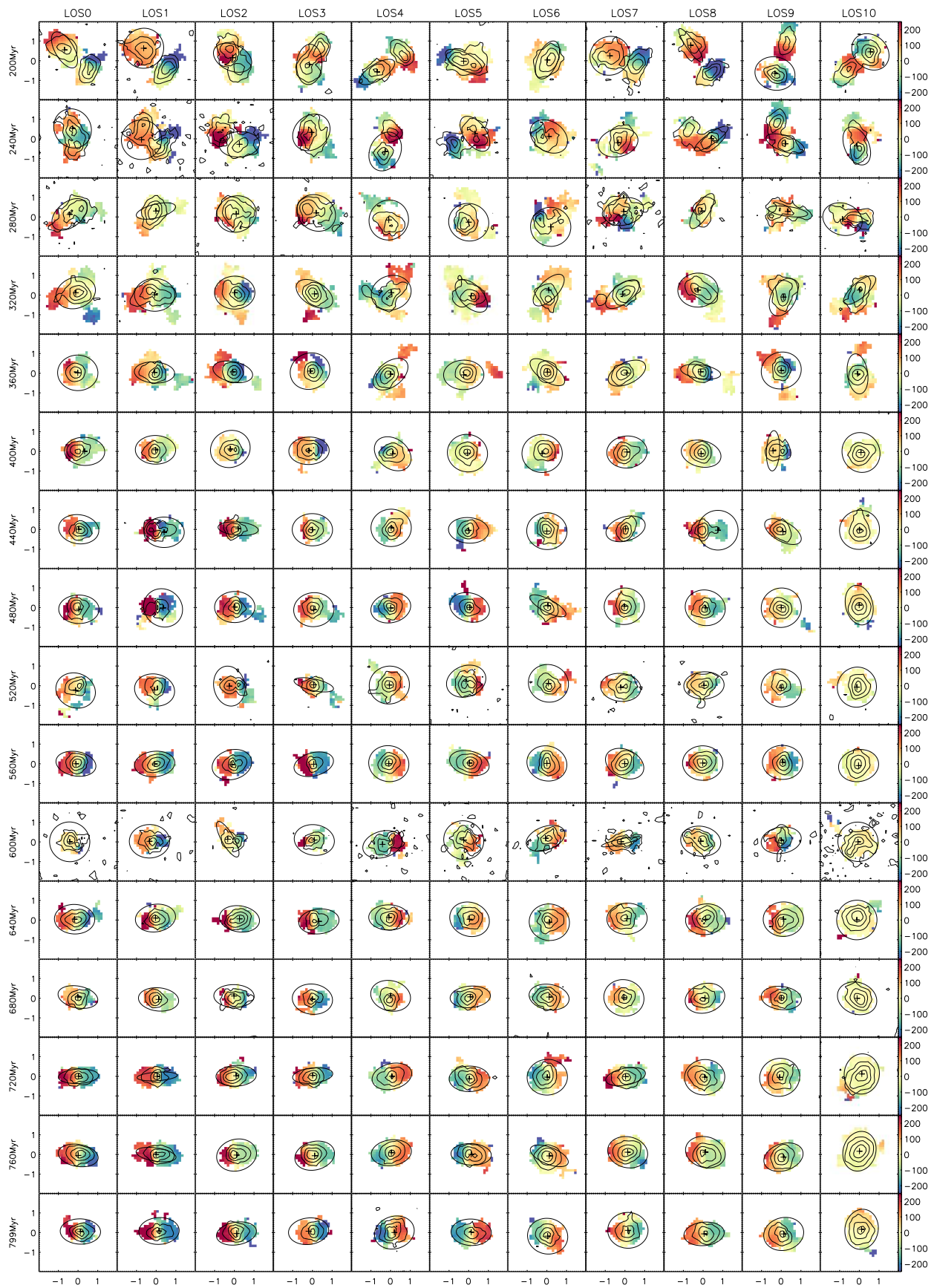
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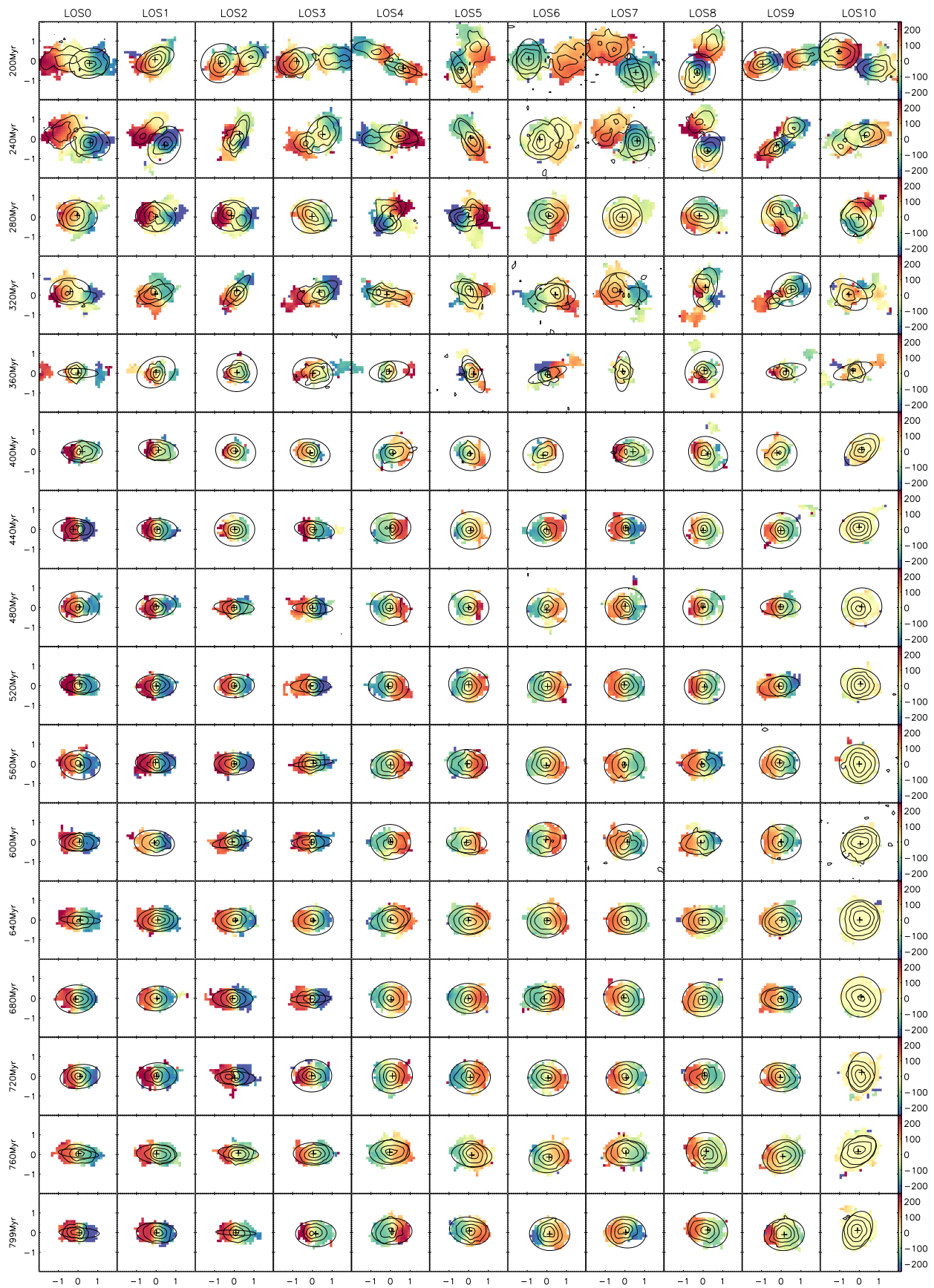
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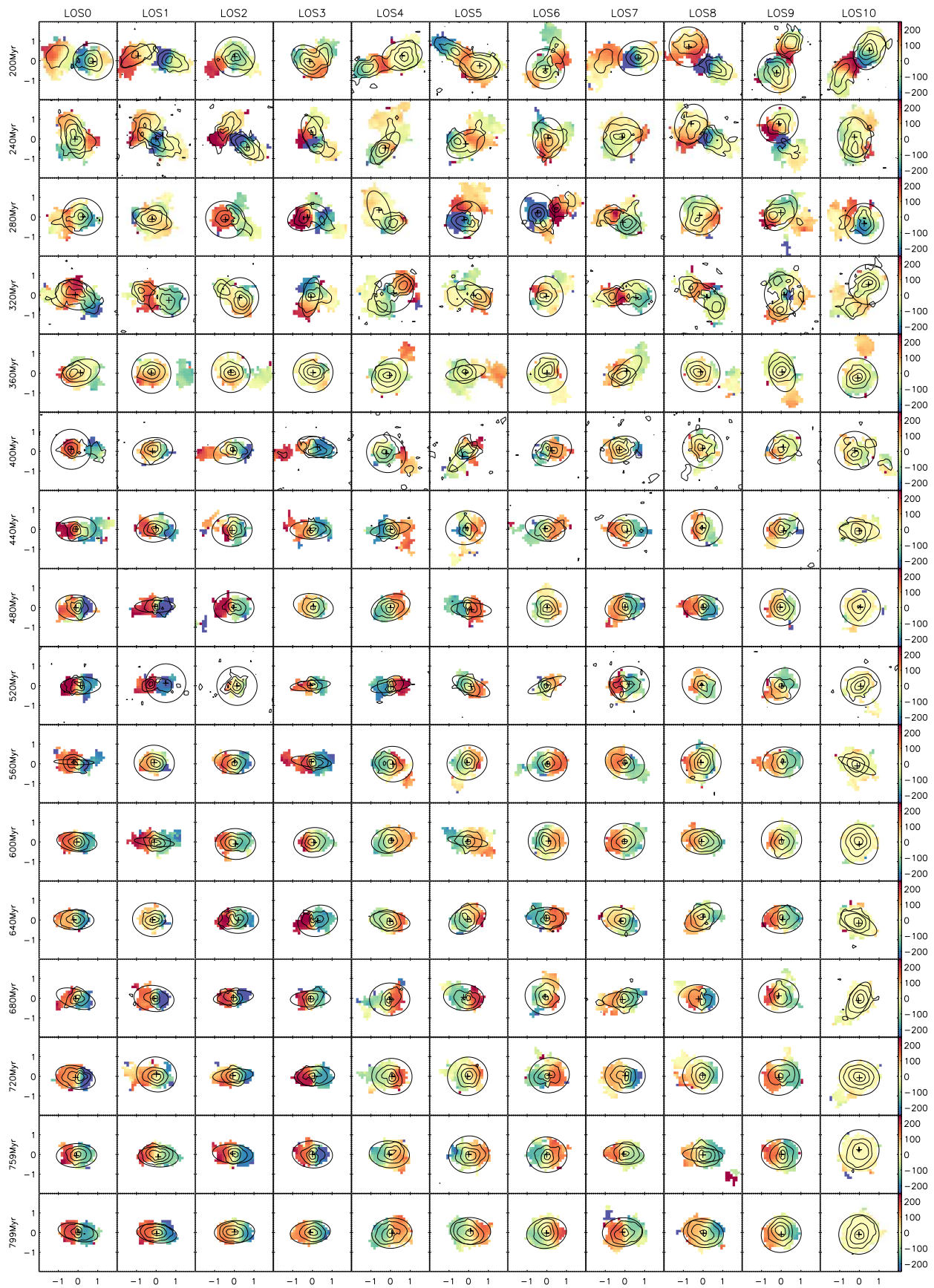
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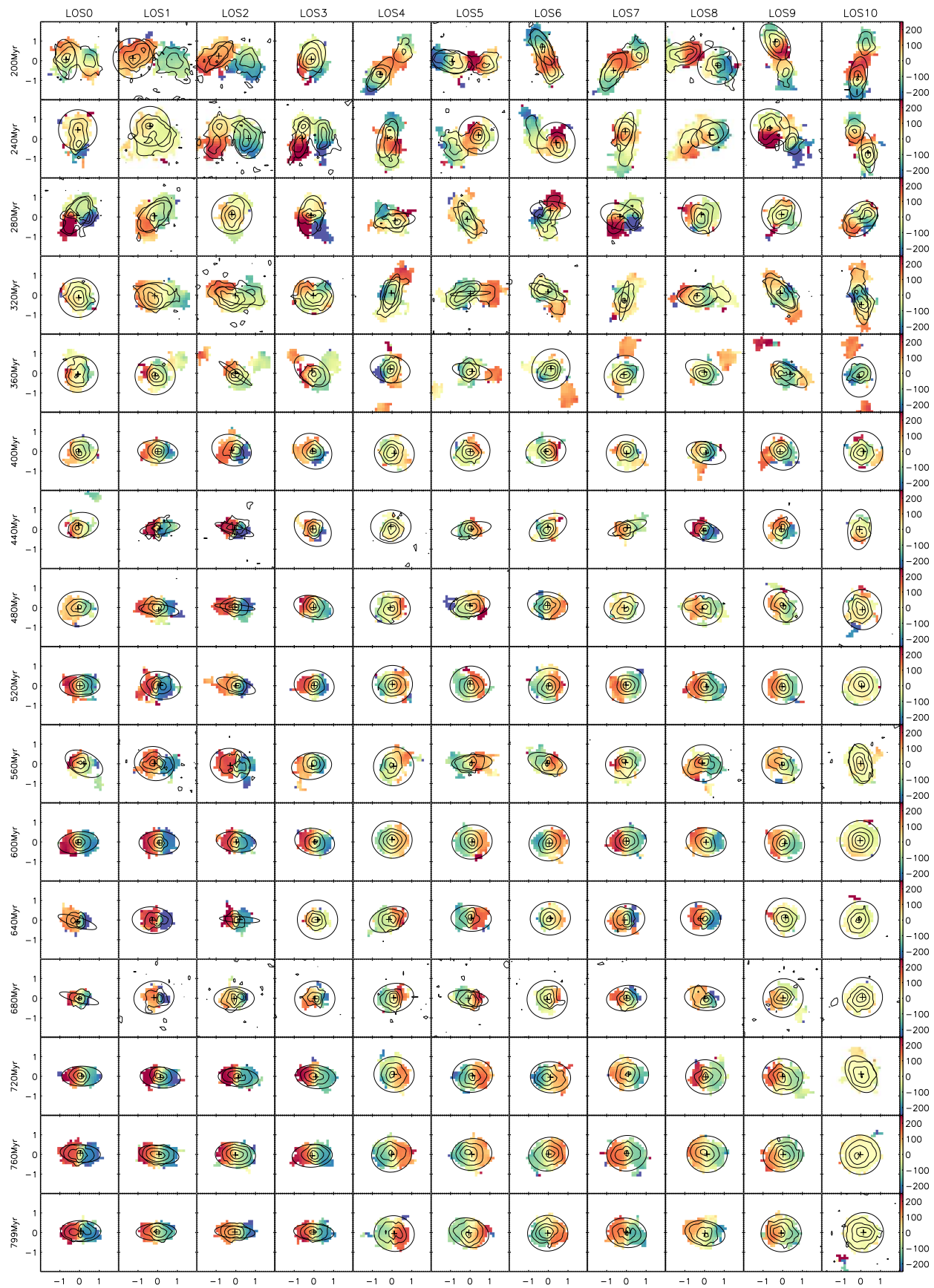
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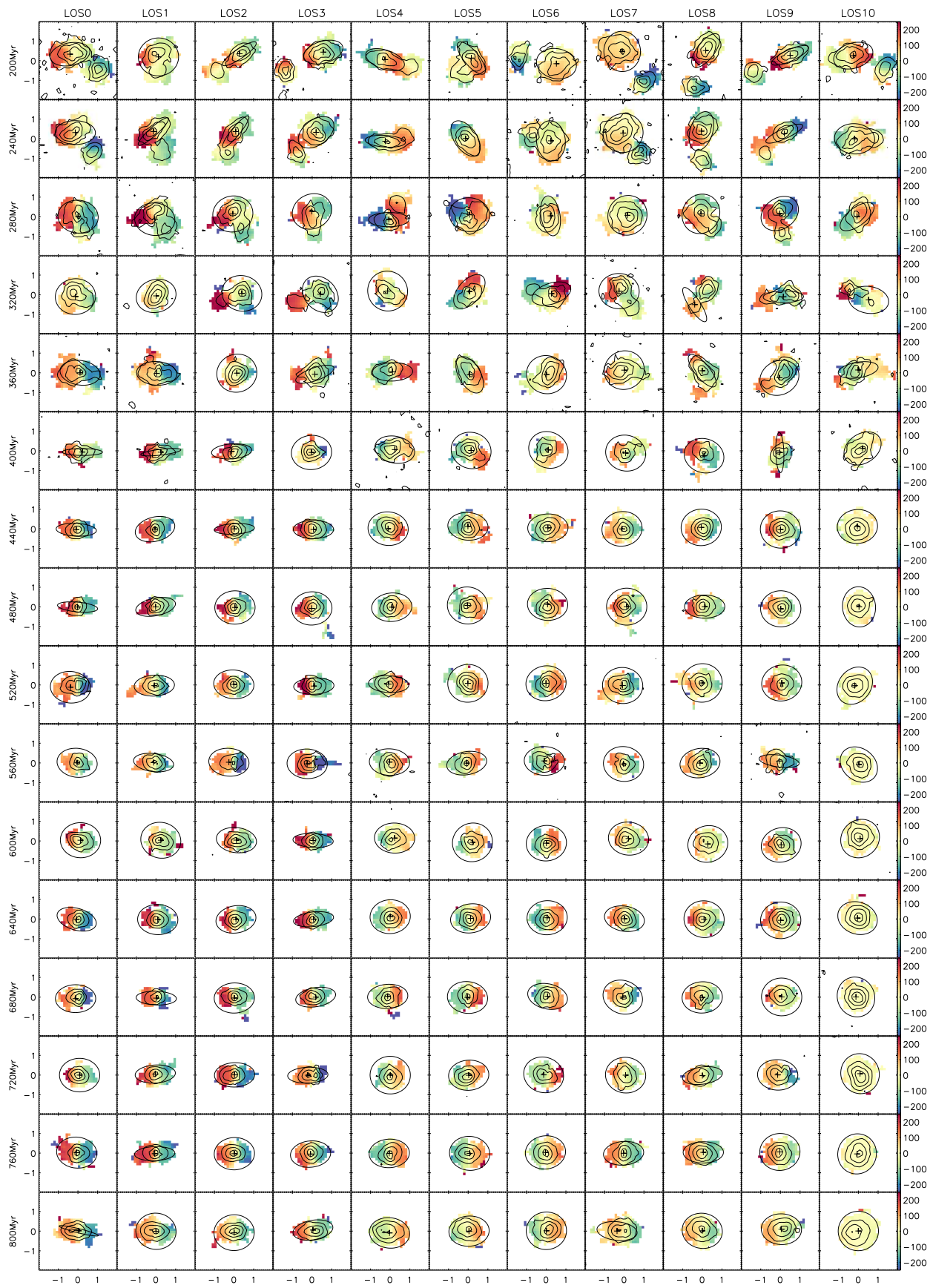
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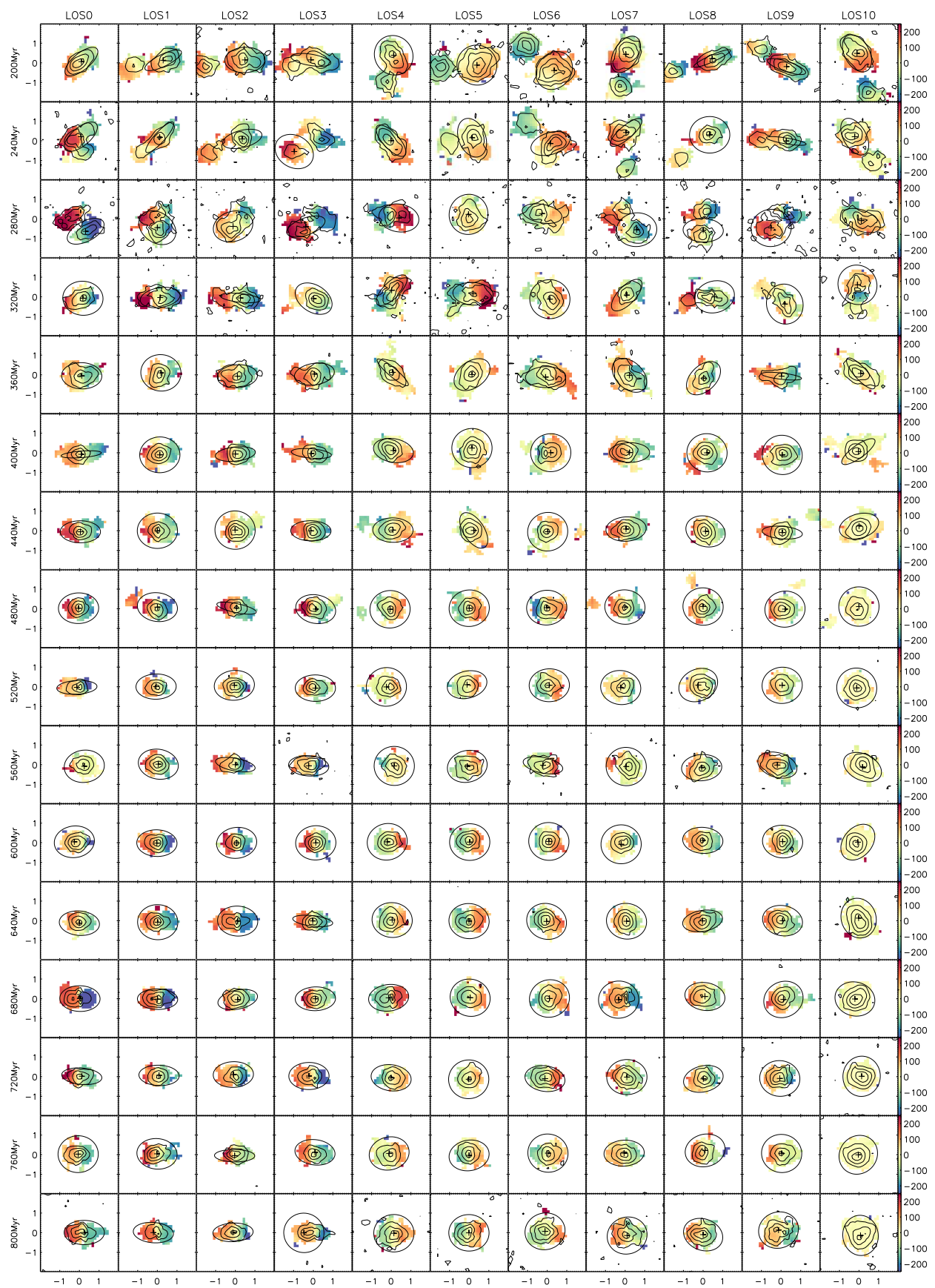
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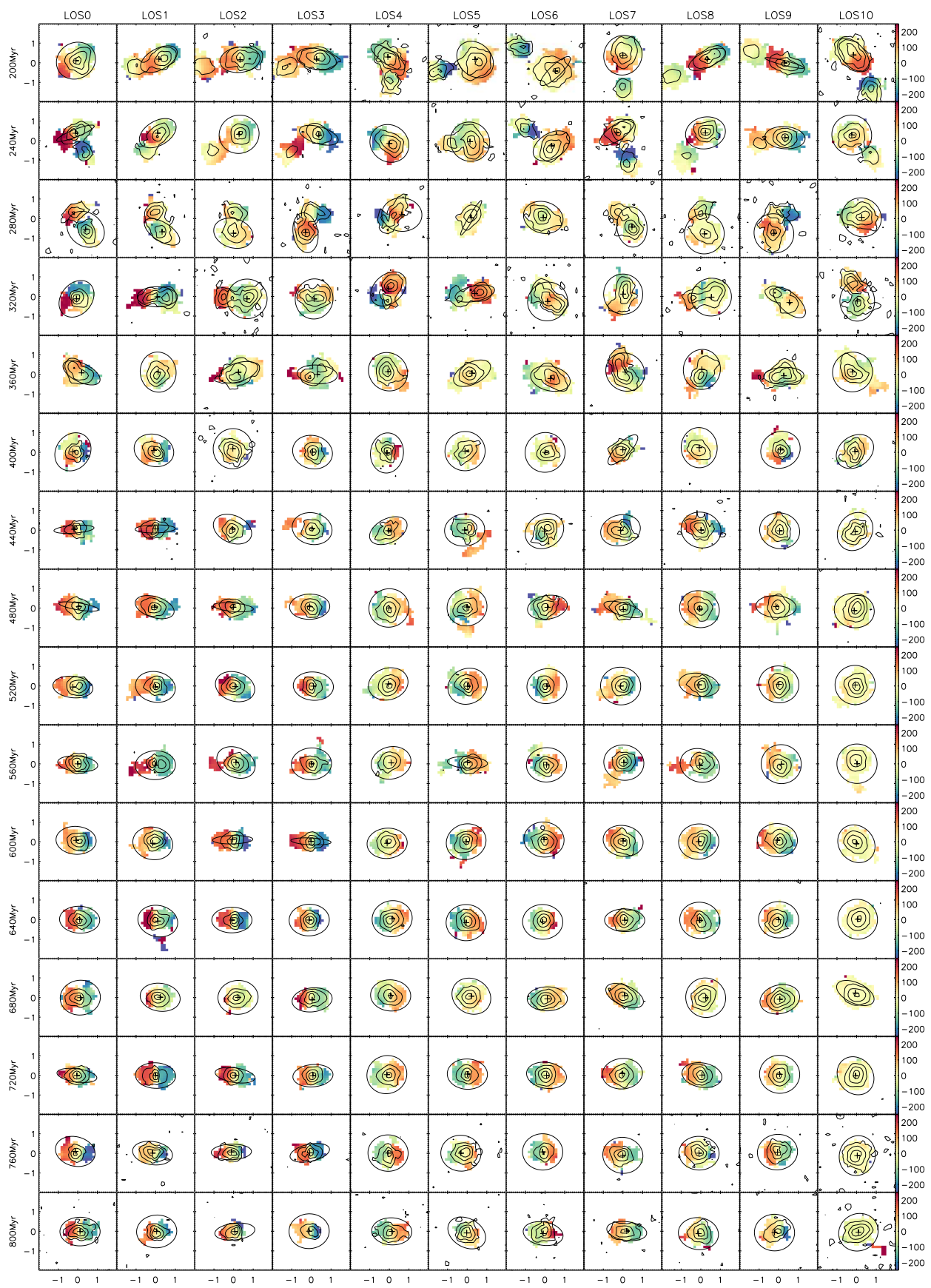
Suite A.3 : simulation de fusion G2_G2_90_90_180.



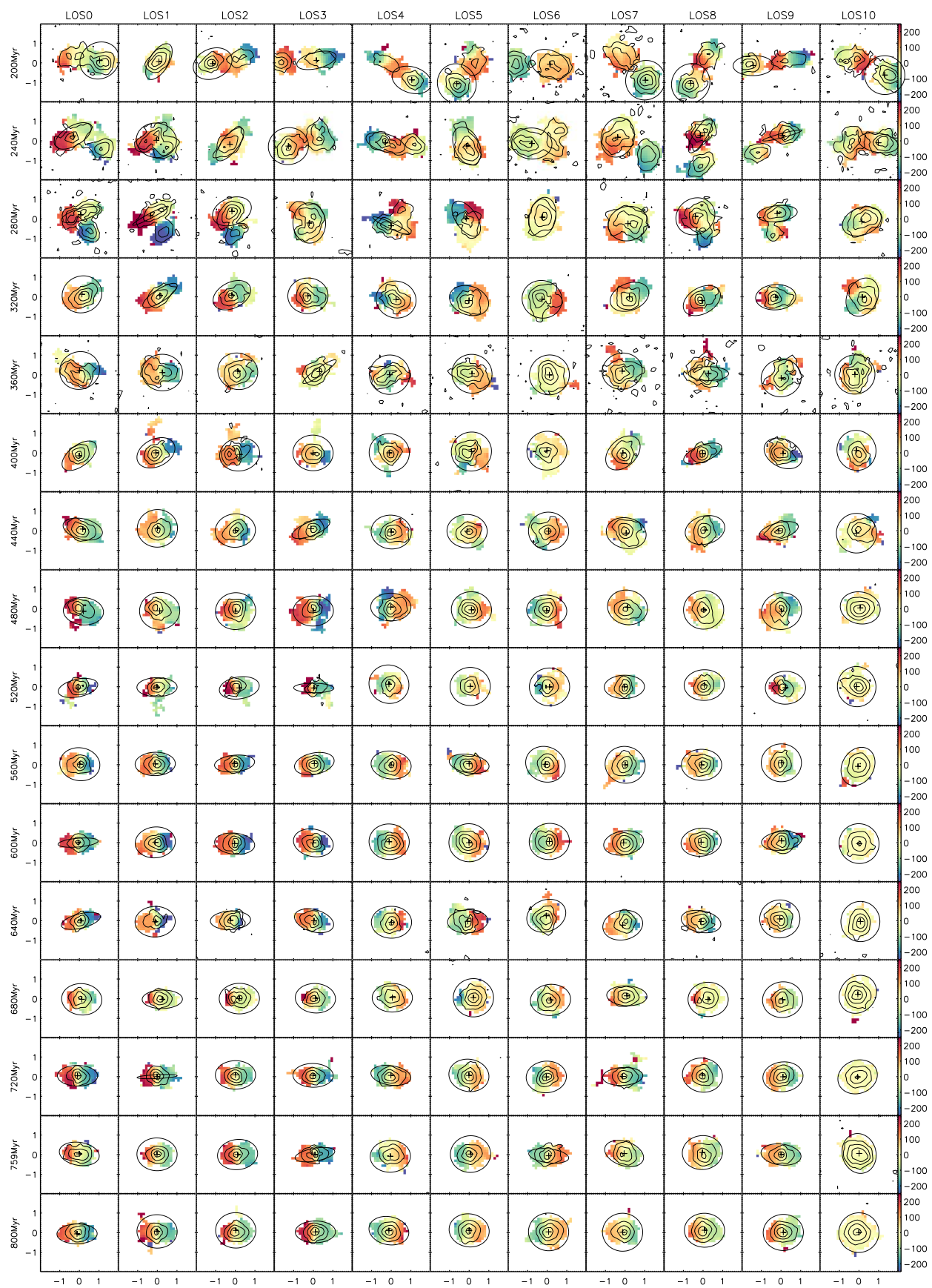
Suite A.3 : simulation de fusion G2_G3_90_90_90.



Suite A.3 : simulation de fusion G2_G3_90_90_0.



Suite A.3 : simulation de fusion G2_G3_90_0_90.



Suite A.3 : simulation de fusion G2_G3_90_90_180.

Articles concernant l'analyse de l'échantillon MASSIV

Résumé

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B.2 Article IV : MASSIV V. The major merger rate of star-forming galaxies at $0.9 < z < 1.8$ from IFS-based close pairs	313

Dans cette annexe, je présente les deux articles auxquels j'ai pu contribuer au cours de cette thèse via mon implication dans le projet MASSIV qui m'a amené à participer à la réduction et l'analyse des données de l'instrument SINFONI. Les deux articles présentés dans cette annexe sont principalement axés sur une analyse cinématique et morphologique des données galaxies MASSIV.

B.1 Article III : MASSIV II. Kinematics and close environment classification

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**Astronomy
&
Astrophysics**

MASSIV: Mass Assembly Survey with SINFONI in VVDS

II. Kinematics and close environment classification^{★,★★,★★★}

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ABSTRACT

Context. Processes driving mass assembly are expected to evolve on different timescales along cosmic time. A transition might happen around $z \sim 1$ as the cosmic star formation rate starts its decrease.

Aims. We aim to identify the dynamical nature of galaxies in a representative sample to be able to infer and compare the mass assembly mechanisms across cosmic time.

Methods. We present an analysis of the kinematics properties of 50 galaxies with redshifts $0.9 < z < 1.6$ from the MASSIV sample observed with SINFONI/VLT with a mass range from $4.5 \times 10^9 M_{\odot}$ to $1.7 \times 10^{11} M_{\odot}$ and a star formation rate from $6 M_{\odot} \text{ yr}^{-1}$ to $300 M_{\odot} \text{ yr}^{-1}$. This is the largest sample with 2D kinematics in this redshift range. We provide a classification based on kinematics as well as on close galaxy environment.

Results. We find that a significant fraction of galaxies in our sample (29%) experience merging or have close companions that may be gravitationally linked. This places a lower limit on the fraction of interacting galaxies because ongoing mergers are probably also present but harder to identify. We find that at least 44% of the galaxies in our sample display ordered rotation, whereas at least 35% are non-rotating objects. All rotators except one are compatible with rotation-dominated ($V_{\text{max}}/\sigma > 1$) systems. Non-rotating objects are mainly small objects ($R_e < 4 \text{ kpc}$). They show an anti-correlation of their velocity dispersion and their effective radius. These low-mass objects ($\log M_{\text{star}} < 10.5$) may be ongoing mergers in a transient state, galaxies with only one unresolved star-forming region, galaxies with an unstable gaseous phase or, less probably, spheroids. Combining our sample with other 3D-spectroscopy samples, we find that the local velocity dispersion of the ionized gas component decreases continuously from $z \sim 3$ to $z = 0$. The proportion of disks also seems to be increasing in star-forming galaxies when the redshift decreases. The number of interacting galaxies seems to be at a maximum at $z \sim 1.2$.

Conclusions. These results draw a picture in which cold gas accretion may still be efficient at $z \sim 1.2$ but in which mergers may play a much more significant role at $z \sim 1.2$ than at higher redshift. From a dynamical point of view, the redshift range $1 < z < 2$ therefore appears as a transition period in the galaxy mass assembly process^{****}.

Key words. galaxies: evolution – galaxies: formation – galaxies: kinematics and dynamics – galaxies: high-redshift

* Based on observations collected at the European Southern Observatory (ESO) Very Large Telescope, Paranal, Chile, as part of the Programs 179.A-0823, 177.A-0837, 78.A-0177, 75.A-0318, and 70.A-9007.

** Based on observations obtained with MegaPrime/MegaCam, a joint project of CFHT and CEA/DAPNIA, at the Canada-France-Hawaii Telescope (CFHT) which is operated by the National Research Council (NRC) of Canada, the Institut National des Sciences de l'Univers of the Centre National de la Recherche Scientifique (CNRS) of France, and the University of Hawaii. This work is based in part on data products produced at TERAPIX and the Canadian Astronomy Data Centre as part of the Canada-France-Hawaii Telescope Legacy Survey, a collaborative project of NRC and CNRS.

*** Appendices are available in electronic form at <http://www.aanda.org>

**** All the data published in this paper are publicly available at the time of the publication following this link: <http://cosmosdb.lambrate.inaf.it/VVDS-SINFONI>.

1. Introduction

During the last decade the first observations of distant galaxies with integral field unit spectrographs (IFU) have led to the construction of several galaxy samples from $z \sim 0.5$ to $z \sim 3$. These IFU observations have given new insights into the resolved physical properties of galaxy populations at various redshifts but a clear view of the transition epoch between young, unstable and clumpy galaxies and evolved and stable galaxies that form the Hubble sequence is still missing.

In the redshift range $0.5 < z < 3$, large quantities of gas are converted into stars, producing a peak in the cosmic star formation at these epochs (see e.g. Hopkins & Beacom 2006; Tresse et al. 2007). Large gas reservoirs are being accreted onto galaxies via various mechanisms: isolated events such as galaxy major and minor mergers (e.g. de Ravel et al. 2009; Conselice et al. 2008; Lin et al. 2008; López-Sanjuan et al. 2011) and cold

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gas accretion along cosmic filaments, a more continuous process (e.g. Kereš et al. 2005; Dekel & Birnboim 2006; Genel et al. 2008; Dekel et al. 2009). Cosmological numerical simulations have helped to outline the importance that these processes may play in the build-up of the present-day Hubble sequence. Indeed, on the one hand, merging galaxies are a natural mass assembly mechanism expected in the Λ CDM framework where dark matter halos grow from hierarchical assembly. Some authors have shown that a peak of major merger activity might occur around $1 < z < 2$ (Ryan et al. 2008; Conselice et al. 2008; López-Sanjuan et al. 2009). On the other hand, cold gas accretion may play a significant role at $z > 2$ but might be less usual at $z < 1$ (e.g. Dekel et al. 2009; Kereš et al. 2009). This process seems to be efficient for halo masses larger than $\sim 10^{11} M_{\odot}$ (Bouché et al. 2010).

Recent IFU studies aim at understanding the role of these different processes in the precursors of local ellipticals and spirals. From the IMAGES sample at $0.4 < z < 0.75$ (Yang et al. 2008; Neichel et al. 2008; Puech et al. 2008; Rodrigues et al. 2008), it has been shown that regular rotating disks are quite similar to local rotators and that merging is playing a significant role in galaxy mass assembly. At higher redshift, the SINS sample at $z \sim 2.2$ (Genzel et al. 2008; Shapiro et al. 2008; Förster Schreiber et al. 2009; Shapiro et al. 2009; Cresci et al. 2009), the LSD/AMAZE sample at $z \sim 3.3$ (Gnerucci et al. 2011) and the sample built by Law et al. (2009) at $2 < z < 3$ are all containing many galaxies with high gaseous turbulence. Based on numerical simulation results of unstable gas-rich disks (e.g. Bournaud et al. 2007; Dekel et al. 2009), these authors claim that continuous cold gas accretion along cosmic web filaments is playing a significant role in mass assembly at these redshifts because they find that half the disks are dispersion-dominated and that even rotation-dominated disks show a high gaseous local velocity dispersion (above 60 km s^{-1}), uncommon for local galaxies (Epinat et al. 2010). Lehnert et al. (2009) and Le Tiran et al. (2011) suggested that this high gaseous local velocity dispersion is not directly powered by shocks or Jeans instabilities due to cold gas accretion but might be related to a vigorous star formation at both high- and low-redshift. This was also supported by Green et al. (2010) who observed a few local analogs to these high-redshift galaxies that all have a star formation rate above $15 M_{\odot}/\text{yr}$. Gonçalves et al. (2010) also observed galaxies with high gaseous local velocity dispersion among a sample of $z \sim 0.2$ Lyman-break analogs that are also forming stars very efficiently. It has been suggested that this intense star formation could also be triggered by merging (Basu-Zych et al. 2007, 2009a,b; Overzier et al. 2008).

Among these various samples (IMAGES, SINS and LSD/AMAZE), the redshift range $0.9 < z < 1.8$, at the peak of the cosmic star formation history, is still poorly explored. The MASSIV survey has been built to study this redshift range, which seems to correspond to the period where the modern Hubble sequence is being built (e.g. Bell et al. 2004; Williams et al. 2009).

The first goal of this paper is to study the dynamical nature of 50 galaxies with $0.9 < z < 1.6$ from MASSIV and the evolution with redshift of the fraction of galaxies in a given dynamical state. The second objective is to identify the main physical mechanisms responsible for the transition between $z \sim 2$, where many disks show high intrinsic gaseous velocity dispersions, and $z \sim 0.5$ where the disks seem to be more “stable”. The general presentation of the MASSIV sample selection and data acquisition strategy is the subject of a companion paper (Contini et al. 2012). An analysis focused on the dynamical properties

of rotators is presented in a second companion paper (Vergani et al. 2012) and the analysis of the spatially-resolved metallicity of this first set of 50 MASSIV galaxies is discussed in a third paper (Queyrel et al. 2012).

The paper is organized as follows. In Sect. 2, the MASSIV “first epoch” sample, the observations, and the data reduction are presented. In Sect. 3, galaxy morphology and kinematics models are described. The galaxy classification scheme is detailed in Sect. 4 and the discussion of this classification is given in Sect. 5. Appendix A contains detailed informations and comments on each galaxy.

In this paper, we use the cosmological parameters $\Omega_m = 0.3$, $\Omega_{\Lambda} = 0.7$ and $H_0 = 70 \text{ km s}^{-1} \text{ Mpc}^{-1}$.

2. Observations and data reduction

2.1. Sample selection

A complete description of the MASSIV sample selection and general properties is presented in Contini et al. (2012). Here we only present a brief summary.

The MASSIV sample contains 84 star-forming galaxies in the redshift range $0.9 < z < 1.8$. It was constructed from the VVDS (VIMOS VLT Deep Survey) spectroscopic survey (Le Fèvre et al. 2005). This survey is *I*-band magnitude limited and complete up to magnitude $I_{AB} < 24$ for VVDS-Deep (RA = 02 h, Le Fèvre et al. 2005), $I_{AB} < 24.75$ for VVDS-Ultra-Deep (RA = 02 h, Cassata et al. 2011; Le Fèvre et al., in prep.) and $I_{AB} < 22.5$ for VVDS-Wide (RA = 14 h and RA = 22 h, Garilli et al. 2008). VVDS provides low-resolution ($R \sim 230$) spectra that lead to accurate spectroscopic redshifts for 4446 galaxies in the redshift range $0.9 < z < 2$.

The selection of the MASSIV sample was based on star formation activity traced by [O II] $\lambda 3727$ emission line equivalent width for galaxies with $z < 1.46$ and from their observed photometric *UBVR*IK spectral energy distribution for galaxies with $z > 1.46$. These criteria ensure that the selected targets are star-forming galaxies for which strong emission lines can be studied using SINFONI to trace the kinematics. The [O II] $\lambda 3727$ selection criteria has been tested on a pilot sample (Epinat et al. 2009) and has proven to be very efficient. Resulting from this selection function, the MASSIV sample provides a good representation of star-forming galaxies with $SFR \geq 5 M_{\odot} \text{ yr}^{-1}$ at $z \sim 1.5$ in the stellar mass regime $10^9 - 10^{11} M_{\odot}$ (see Contini et al. 2012 for a detailed study).

The MASSIV “first epoch” sample discussed in this paper is a subsample of the MASSIV sample. It contains the 50 galaxies with $0.9 < z < 1.6$ (the median redshift is 1.24) that were observed before January 2010. Except for two galaxies observed in the VVDS-Ultra-Deep, the galaxies studied in this paper are from the VVDS-Deep and VVDS-Wide samples.

2.2. Observations

SINFONI (Eisenhauer et al. 2003; Bonnet et al. 2004) was used to obtain 2D spatially resolved spectroscopy in the NIR of the MASSIV galaxies around the $H\alpha$ line or around the [O III] $\lambda\lambda 4959, 5007$ line for four galaxies. Out of an initial sample of 50 galaxies no line was detected for only four galaxies. The success rate is much better when the $H\alpha$ line is targeted (44/46). VVDS220148046 was observed at $z = 2.244$, whereas it was expected at $z \sim 1.371$ from the VIMOS spectrum. The redshift determination for this galaxy is based on the observation of [O III] $\lambda\lambda 4959, 5007$ and $H\beta$ lines in our SINFONI data. The observations were obtained in service mode from period

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P79 to period P82 as part of the Large Programme 179.A-0823 (P.I. T. Contini) and during a pilot program presented in Epinat et al. (2009) (ESO runs 75.A-0318 and 78.A-0177). Galaxies were observed in two or three observing blocks.

For the galaxies with $z < 1.1$, $H\alpha$ was observed with the J grism over the spectral range 1.08–1.41 μm , whereas for galaxies with $z > 1.2$, $H\alpha$ was observed with the H grism over the spectral range 1.43–1.86 μm . $[\text{O III}] \lambda\lambda 4959, 5007$ was observed in the J -band except for VVDS220148046 for which it was observed in the H -band. SINFONI spectral resolution in J - and H -bands reaches ~ 2000 and ~ 2500 , respectively.

For seeing-limited observations, we offset the target alternately from one corner of the field-of-view of the instrument to the opposite one (object nodding) to maximize the observing efficiency during the nights. This observing strategy allowed us to avoid sky frame acquisition. For AO observations, we used one sky frame for four object frames to maximize the time spent on sources (details in Contini et al. 2012). In addition, we also applied a sub-dithering to avoid the return of the target to the same position on the chip. To allow for an accurate on-source pointing of our galaxies, we acquired them through a blind offset from a bright nearby star (PSF star used to measure the spatial PSF) to our target. We also observed standard (STD) stars for flux calibration during the same night. Individual exposures were 300 s, 600 s or 900 s with a total on-source integration time that ranges between 1 h and 2 h.

Most of the data were observed in seeing-limited mode using the $0.125'' \times 0.25''$ pixel scale leading to a $8'' \times 8''$ field of view with a mean seeing of $0.68 \pm 0.12''$, considering only detected galaxies. However, a subset of seven galaxies were observed with the laser guide star (LGS) adaptive optics system using the $0.05'' \times 0.10''$ pixel scale with a $3.2'' \times 3.2''$ field of view. Two of them were not detected (VVDS020126402 and VVDS220071601) and the PSF star was missed for VVDS220386469. The mean spatial resolution for the four other galaxies is $0.23''$. The observing details along with the periods at which the galaxies were observed are listed in Table 1.

2.3. Data reduction

The data reduction was performed using the ESO-SINFONI pipeline (version 2.0.0, Modigliani et al. 2007) complemented with additional IDL and PYTHON routines to perform the data processing homogeneously among the reducers and to improve some reduction steps.

First, the PSF stars and STD stars observations were reduced using standard data reduction.

Then, for each object science frame, the following steps were applied:

1. bad line removal in the raw data;
2. dark-current, sky-background and night-sky line subtraction from the raw data using the contiguous frame (with the target in the opposite corner due to the observing strategy or with only sky for AO data);
3. flat-field correction using an internal lamp;
4. wavelength calibration using arc-lamps;
5. flux calibration using the STD telluric star spectrum with standard techniques to convert counts into flux units. This enables one to correct for the atmospheric transmission and instrumental response;
6. cube reconstruction (with and without sky lines) in counts;

7. adjustment of the astrometry to match our I -band reference images (CFHT12k survey, McCracken et al. 2003) using the acquisition PSF star and the offsets of object observations with respect to this star. This method was not possible on AO data due to missing meta-informations. For some galaxies of the pilot program, no PSF star was observed and in other cases, PSF stars were saturated in the CFHT images, which provided a deteriorated astrometry.

In the pipeline, the spaxels are resampled so that they are square ($0.125''$ or $0.05''$).

All cubes obtained for each science frame for a given object were then combined using the SINFO_UTL_CUBE_COMBINE recipe to obtain the final cubes using an additional median filtering to remove sky line residuals as accurately as possible from the final cube.

Mono-dimensional sky spectra were extracted from the sky cubes. They were used to

- determine the effective spectral resolution: using a Gaussian fit to approximate sky lines, the spectral resolution element is found to be fairly constant over the whole wavelength range in the J - and H -bands and the dispersion of the Gaussian was estimated to $\sigma \sim 2.8 \pm 0.2 \text{ \AA}$;
- quantify the noise that was considered as a Poissonian noise.

The data from the pilot study (Epinat et al. 2009) were reduced again using this new procedure.

2.4. Map extraction

The ionized gas kinematics of MASSIV galaxies is studied through the brightest emission line available in the NIR spectra, the $H\alpha$ line or the $[\text{O III}] \lambda 5007$ line in a few cases. IDL routines, based on the `mpfit` routine (Markwardt 2009), were used to extract the kinematic maps from the SINFONI data. First, to enhance the signal-to-noise ratio (S/N) without degrading the spatial resolution of the data, a subresolution 2D spatial Gaussian smoothing (FWHM of two pixels) was applied on the data cubes. In addition, the residual cosmic rays were rejected using a 15σ clipping on 3×3 pixel boxes. Then, for each spatial pixel, the spectrum around $H\alpha$ was fitted by a single Gaussian profile and a constant continuum (four parameters in total). To minimize the effects of noise induced by sky lines but also of sky line residuals on the line parameter determination, the 1D sky spectrum was used as an estimate of the noise to weight the contribution of each spectral element. From these fitting techniques it was possible to recover the line flux map, the velocity field and the velocity dispersion map for each source. The instrumental spectral PSF was taken into account to compute the velocity dispersion: $\sigma^2 = \sigma_{\text{obs}}^2 - \sigma_{\text{PSF}}^2$. During the line fitting procedure, the velocity dispersion was allowed to vary in the range 40 km s^{-1} (to avoid fitting noise) to 250 km s^{-1} (to avoid fitting a continuum). 2D error maps were also derived for each quantity from the fitting procedure. These are statistical errors that take into account the error spectrum and that indicate the accuracy of the fit for each parameter. An S/N map was computed. The computation of this map was refined since the study of the pilot program (Epinat et al. 2009). The presence of sky lines is now taken into account to modulate the confidence on the line detection: the inverse of the sky spectrum is used for weighting (w). The signal (S) is computed as the weighted flux of the line divided by the dispersion of the line (σ_λ) multiplied by $\sqrt{2\pi}$. In the following equations, the weight was normalized ($\sum w = 1$), n_z is the number of spectral elements, $\Delta\lambda$ refers to the spectral sampling and

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Table 1. MASSIV galaxies “first epoch” sample and SINFONI observation setups.

VVDS ID	RA	Dec	z	Scale	Pixel scale	Band	Line	t_{exp}	Seeing	R	Period
(1)	[J2000]	[J2000]	(4)	[kpc $''$]	[$''$]	(7)	(8)	[min]	[$''$]	(11)	(12)
020106882	02:25:21.819	-04:46:18.35	1.3991	8.43	0.125 × 0.25	<i>H</i>	H α	80	0.49	2477	P82G
020116027	02:25:51.085	-04:45:06.08	1.5302	8.47	0.125 × 0.25	<i>H</i>	H α	75	0.60	2609	P75A
020126402 ¹	02:25:11.658	-04:43:40.12	1.2332	8.33	0.050 × 0.10	<i>J</i>	[OIII]	60	–	–	P79B
020147106	02:26:45.362	-04:40:47.50	1.5195	8.47	0.125 × 0.25	<i>H</i>	H α	120	0.65	2598	P75A
020149061	02:27:05.226	-04:40:29.21	1.2905	8.37	0.125 × 0.25	<i>H</i>	H α	80	0.85	2362	P82G
020164388	02:26:50.942	-04:38:20.72	1.3547	8.41	0.125 × 0.25	<i>H</i>	H α	80	0.83	2430	P82G
020167131	02:26:47.307	-04:37:55.36	1.2246	8.32	0.125 × 0.25	<i>J</i>	[OIII]	120	0.68	2295	P79B
020182331	02:26:44.242	-04:35:52.01	1.2290	8.32	0.125 × 0.25	<i>H</i>	H α	180	0.74	2302	P78A
020193070	02:25:18.713	-04:34:19.77	1.0279	8.06	0.125 × 0.25	<i>J</i>	H α	120	0.58	2094	P82G
020208482	02:25:16.739	-04:32:11.92	1.0375	8.08	0.125 × 0.25	<i>J</i>	H α	120	0.58	2104	P82G
020214655	02:26:23.441	-04:31:22.78	1.0395	8.08	0.125 × 0.25	<i>J</i>	H α	80	0.87	2101	P82G
020217890 ¹	02:26:27.162	-04:30:51.83	1.5129	8.46	0.125 × 0.25	<i>H</i>	H α	120	–	–	P78A
020239133	02:26:43.006	-04:28:31.20	1.0194	8.04	0.125 × 0.25	<i>J</i>	H α	80	0.79	2084	P82G
020240675	02:26:54.140	-04:28:17.64	1.3270	8.40	0.125 × 0.25	<i>H</i>	H α	80	0.85	2402	P82G
020255799	02:26:45.859	-04:26:15.80	1.0351	8.07	0.125 × 0.25	<i>J</i>	H α	80	0.76	2101	P82G
020261328	02:27:11.023	-04:25:31.57	1.5290	8.47	0.125 × 0.25	<i>H</i>	H α	60	0.62	2609	P75A
020278667	02:25:58.203	-04:23:11.67	1.0516	8.10	0.125 × 0.25	<i>J</i>	H α	120	0.65	2115	P82G
020283083	02:26:30.832	-04:22:35.82	1.2818	8.36	0.125 × 0.25	<i>H</i>	H α	80	0.78	2353	P82G
020283830	02:26:28.926	-04:22:31.14	1.3949	8.43	0.125 × 0.25	<i>H</i>	H α	120	0.77	2472	P82G
020294045	02:25:47.152	-04:21:07.41	1.0028	8.01	0.125 × 0.25	<i>J</i>	H α	120	0.59	2067	P79B
020306817 ¹	02:25:50.316	-04:19:22.93	1.2225	8.32	0.125 × 0.25	<i>J</i>	[OIII]	120	–	–	P79B
020363717	02:26:23.709	-04:11:57.87	1.3339	8.40	0.125 × 0.25	<i>H</i>	H α	80	0.64	2407	P82G
020370467	02:26:14.690	-04:11:05.44	1.3338	8.40	0.125 × 0.25	<i>H</i>	H α	80	0.71	2407	P82G
020386743	02:27:13.989	-04:08:59.73	1.0487	8.09	0.125 × 0.25	<i>J</i>	H α	120	0.73	2111	P79B
020461235	02:26:47.102	-04:23:55.70	1.0349	8.07	0.125 × 0.25	<i>J</i>	H α	120	0.63	2101	P79B
020461893	02:27:12.252	-04:23:11.28	1.0486	8.09	0.125 × 0.25	<i>J</i>	H α	80	0.60	2115	P82G
020465775	02:26:59.366	-04:19:00.08	1.3583	8.41	0.125 × 0.25	<i>H</i>	H α	80	0.88	2434	P82G
140083410	13:57:50.595	+04:17:38.71	0.9435	7.89	0.125 × 0.25	<i>J</i>	H α	80	0.69	2005	P81D
140096645	13:58:26.336	+04:19:47.75	0.9655	7.94	0.125 × 0.25	<i>J</i>	H α	120	0.56	2024	P81D
140123568	13:55:57.628	+04:24:20.11	1.0012	8.01	0.125 × 0.25	<i>J</i>	H α	120	0.76	2067	P79A
140137235	13:56:12.729	+04:26:31.74	1.0445	8.09	0.125 × 0.25	<i>J</i>	H α	80	0.76	2111	P79A
140217425	13:57:56.405	+04:38:37.00	0.9792	7.97	0.125 × 0.25	<i>J</i>	H α	100	0.95	2040	P81D
140258511	14:00:19.658	+04:44:45.86	1.2423	8.33	0.125 × 0.25	<i>H</i>	H α	80	0.49	2315	P79A
140262766	13:59:55.518	+04:45:30.04	1.2836	8.37	0.125 × 0.25	<i>H</i>	H α	120	0.51	2353	P79A
140545062	13:59:35.598	+05:30:31.11	1.0408	8.08	0.125 × 0.25	<i>J</i>	H α	120	0.70	2104	P81D
220014252	22:17:45.677	+00:28:39.52	1.3105	8.38	0.125 × 0.25	<i>H</i>	H α	120	0.70	2384	P75A
220015726	22:15:42.435	+00:29:03.58	1.2933	8.37	0.125 × 0.25	<i>H</i>	H α	120	0.46	2384	P79B
220071601 ¹	22:18:01.569	+00:45:34.69	1.3538	8.41	0.050 × 0.10	<i>H</i>	H α	80	–	–	P79B
220148046	22:14:37.904	+01:08:20.65	2.2442	8.24	0.050 × 0.10	<i>H</i>	[OIII]	80	0.27	2450	P81E
220376206	22:20:05.772	-00:08:21.74	1.2445	8.34	0.125 × 0.25	<i>H</i>	H α	120	0.50	2315	P79B
220386469	22:19:56.603	-00:03:03.78	1.0226	8.05	0.050 × 0.10	<i>J</i>	H α	40	0.23 ²	2090	P79B
220397579	22:20:36.512	+00:01:46.85	1.0379	8.08	0.125 × 0.25	<i>J</i>	H α	120	0.64	2101	P79B
220544103	22:15:25.689	+00:06:40.31	1.3973	8.43	0.125 × 0.25	<i>H</i>	H α	120	0.76	2472	P75A
220544394	22:14:24.153	+00:06:46.67	1.0101	8.03	0.125 × 0.25	<i>J</i>	H α	120	0.58	2073	P79B
220576226	22:16:11.417	+00:16:30.46	1.0217	8.05	0.125 × 0.25	<i>J</i>	H α	120	0.58	2087	P79B
220578040	22:17:04.113	+00:16:56.80	1.0462	8.09	0.125 × 0.25	<i>J</i>	H α	120	0.62	2111	P79B
220584167	22:15:22.917	+00:18:48.82	1.4655	8.45	0.125 × 0.25	<i>H</i>	H α	120	0.75	2541	P75A
220596913	22:14:29.179	+00:22:18.93	1.2658	8.35	0.050 × 0.10	<i>H</i>	H α	120	0.18	2340	P79B
910193711	02:25:46.285	-04:32:33.43	1.5564	8.47	0.050 × 0.10	<i>H</i>	H α	80	0.27	2636	P82F
910279515	02:25:36.233	-04:21:16.13	1.4013	8.43	0.050 × 0.10	<i>H</i>	H α	80	0.21	2477	P82F

Notes. The coordinates of the galaxies are given in Cols. (2) and (3). The redshift (4) is determined from SINFONI data. The physical scale (5) is computed from the redshift. The SINFONI pixel scale is given in Col. (6). In Col. (7), the wavelength band used to observe the emission line (8) is given. Column (9) is the on-source exposure time. The seeing of SINFONI observations (10) is derived from observed PSF stars. Column (11) is the spectral resolution. The ESO observing period is given in Col. (12). ⁽¹⁾ These galaxies are those for which no line was detected in SINFONI data. The redshift is the one derived from VIMOS spectra, the line is the targeted one and both spectral and spatial resolution have not been measured. ⁽²⁾ No PSF star was observed: the resolution is the mean resolution of AO observations.

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is expressed in the same unit as σ_λ , m is the modeled spectrum, c is the continuum of the modeled spectrum and l is the observed line spectrum:

$$S = \frac{n_z \Delta \lambda \sum w(m - c)}{\sigma_\lambda \sqrt{2\pi}}.$$

The noise (N) is computed as the square root of the weighted variance of the residual spectrum:

$$N = \sqrt{\sum w(l - m)^2}.$$

This yields the S/N :

$$\frac{S}{N} = \frac{n_z \Delta \lambda \sum w(m - c)}{\sigma_\lambda \sqrt{2\pi \sum w(l - m)^2}}. \quad (1)$$

If the weight is constant, then, according to Eq. (1), the S/N is simply the ratio of the intensity of a Gaussian line over the rms of the spectrum. This criterion is robust and an S/N threshold of three was used to clean the kinematic maps (see Appendix B). Extra-cleaning was performed manually to remove isolated pixels (less than about five adjacent pixels) or groups of pixels associated with data artifacts that could have passed the criteria. The existence of any I -band counterpart was checked when there were groups of more than four pixels.

3. Modeling

3.1. Morphology modeling

The stellar continuum is barely detected in the SINFONI data. To study the morphology of the stellar component, we used the best CFHT I -band images available, i.e. from the CFH12K/CFHT survey (McCracken et al. 2003) for the galaxies in the 14 h field and from the CFHT Legacy Survey¹ with the best seeing for the galaxies in the 02 h and 22 h fields. We ran GALFIT (Peng et al. 2002) on those images using a Sersic profile (Sersic 1968) to recover the morphological parameters: the center, the position angle of the major axis (PA), the axis ratio (b/a), the effective radius (R_e), the Sersic index (n) and the total magnitude. GALFIT convolves the model to the spatial PSF to converge into a set of beam-smearing corrected parameters.

To obtain robust estimates of the parameters and of their associated error bars, it is mandatory to control the PSF as well as possible. Indeed, for objects with sizes comparable to the resolution, the use of an overestimated PSF leads to low axis-ratios with small error bars since no strong elongation may be reproduced: this would lead to systematic underestimating of structures in the galaxies. We are indeed in most cases in this situation when the size on the objects is of the same order as the spatial resolution for seeing-limited ground-based imaging surveys. Thus, in each CFHT field, we randomly selected stars to characterize and follow the variation of the PSF. The detection of stars is based on color and morphological criteria for the 02 h and 22 h fields and on morphological criteria only for the 14 h field. From these stars, the width, the axis ratio and the orientation of the PSF are found to vary much across the 14 h and 22 h fields. The most affected field is the 14 h one where the seeing smoothly varies from $0.6''$ to $1.0''$. Therefore, we selected specific PSFs in each field, located at $45''$ from each galaxy on average. The study of the PSF distribution over the various fields enabled us to determine the uncertainty on the PSF FWHM to be

~ 0.1 pixel (~ 20 mas). It is estimated as the deviation from large-scale variations.

GALFIT produces residual maps that were used to check the convergence of the fits. It was sometimes necessary to fit secondary objects to have a correct fit (cf. Appendix A). The parameters of the fits are summarized in Table 2.

We tried in a second step to fit the morphology while fixing the position angle of the major axis to the value determined from the kinematics modeling (cf. Sect. 3.2) as was done in the pilot study (Epinat et al. 2009). However, this leads to a bias toward high axis ratios because GALFIT cannot match the elongation when the position angle is fixed and finds that round morphologies match the data better. Therefore, we adopt the morphologies as derived in an unconstrained way from the CFHT images in the following.

3.2. Kinematics modeling

Among the various dynamical states of galaxies, the easiest to probe is that of the rotating disk. We therefore tested the likelihood of this hypothesis for the galaxies in our sample and recovered the fundamental dynamical parameters within this hypothesis. The velocity field is accordingly fitted with a model that assumes that the ionized gas is located in an infinitely thin rotating disk, as in Epinat et al. (2009). The rotation curve is described by a linear slope in the inner parts and a plateau in the outer parts. The velocity along the line of sight is computed taking into account geometrical position effects. The model parameters are

- x_c, y_c : the center coordinates;
- z : the redshift corresponding to the systemic velocity;
- i : the inclination of the gaseous disk;
- PA_k : the position angle of the major axis;
- V_t : the plateau rotation velocity;
- r_t : the turnover radius at which the plateau is reached.

The method used to adjust the models is described in detail in Epinat et al. (2010). It is based on a χ^2 minimization and takes the velocity error map into account to minimize the contribution of the regions with low S/N . The spatial PSF is taken into account in these models and is described with a 2D Gaussian. Its FWHM is computed on the PSF stars associated to each observation. To compute the model velocity field, a higher resolution velocity field is constructed (with at least eight pixels in the PSF FWHM) from the analytical model and a high-resolution line flux map has also to be built. Indeed, in a final low spatial resolution element, the contribution of the line of sight velocity at higher resolution is weighted by the true line flux distribution. We used a linear interpolation of the observed flux map to avoid making any assumption on the real flux distribution. This is one of the major uncertainties of our models because the real line flux distribution could be more clumpy than observed.

Our model allows us to compute the circular velocity within the disk hypothesis. This assumption is not realistic for all objects in the MASSIV sample in which we also expect mergers (ongoing or late stage), spheroids, or structures with chaotic motions. However, this hypothesis allows us to compute a map that contains only the beam-smearing effect on the velocity dispersion, due to the blurring of large-scale motions. Thus, by subtracting quadratically the map deduced from the model to the observed velocity dispersion map, a velocity dispersion map corrected for the beam-smearing effect is obtained (noted “ σ residuals” in Appendix B). This correction is also valid at first order for non-rotating objects, even if it can overestimate the correction in

¹ <http://www.cfht.hawaii.edu/Science/CFHLS/>

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Table 2. CFHT observations and morphological parameters from GALFIT modeling.

VVDS ID	Survey	Pixel scale [""]	Seeing [""]	PA [°]	b/a	R_e [kpc]	n
(1)	(2)	(3)	(4)	(5)	(6)	(7)	(8)
020106882	LS	0.186	0.604	294 ± 4	0.62 ± 0.03	3.51 ± 0.14	0.77 ± 0.16
020116027	LS	0.186	0.604	184 ± 2	0.38 ± 0.03	4.27 ± 0.16	0.70 ± 0.15
020126402	LS	0.186	0.615	15 ± 8	0.56 ± 0.10	2.12 ± 0.31	0.59 ± 0.58
020147106	LS	0.186	0.615	310 ± 1	0.19 ± 0.15	1.17 ± 0.41	0.41 ± 1.82
020149061	LS	0.186	0.649	201 ± 369	0.93 ± 0.83	1.09 ± 0.69	0.11 ± 6.68
020164388	LS	0.186	0.618	162 ± 3	0.71 ± 0.02	2.72 ± 0.05	0.85 ± 0.09
020167131	LS	0.186	0.618	272 ± 3	0.60 ± 0.03	2.60 ± 0.09	1.64 ± 0.26
020167131s	LS	0.186	0.618	57 ± 1	0.37 ± 0.02	4.11 ± 0.08	0.31 ± 0.09
020182331	LS	0.186	0.618	268 ± 4	0.49 ± 0.04	4.19 ± 0.17	0.63 ± 0.21
020193070	LS	0.186	0.615	215 ± 2	0.33 ± 0.03	3.49 ± 0.09	0.47 ± 0.13
020208482	LS	0.186	0.600	355 ± 7	0.70 ± 0.05	3.67 ± 0.20	0.82 ± 0.21
020214655	LS	0.186	0.604	32 ± 4	0.41 ± 0.06	1.47 ± 0.08	1.97 ± 0.53
020217890	LS	0.186	0.604	171 ± 2	0.21 ± 0.04	3.57 ± 0.14	1.13 ± 0.22
020239133	LS	0.186	0.608	123 ± 2	0.39 ± 0.03	3.14 ± 0.09	0.25 ± 0.13
020240675	LS	0.186	0.619	190 ± 13	0.62 ± 0.13	1.09 ± 0.16	2.03 ± 1.19
020255799	LS	0.186	0.607	164 ± 12	0.80 ± 0.08	1.89 ± 0.26	0.24 ± 0.44
020261328	LS	0.186	0.634	171 ± 8	0.51 ± 0.11	1.83 ± 0.27	0.64 ± 0.53
020278667	LS	0.186	0.596	140 ± 12	0.74 ± 0.09	2.82 ± 1.66	9.60 ± 5.68
020283083	LS	0.186	0.596	301 ± 3	0.38 ± 0.05	4.29 ± 0.20	0.46 ± 0.18
020283830	LS	0.186	0.596	142 ± 1	0.35 ± 0.01	6.82 ± 0.16	0.09 ± 0.06
020294045	LS	0.186	0.606	1 ± 3	0.61 ± 0.03	2.89 ± 0.10	1.00 ± 0.17
020306817	LS	0.186	0.606	86 ± 2	0.54 ± 0.02	4.22 ± 0.09	0.27 ± 0.08
020363717	LS	0.186	0.626	154 ± 6	0.55 ± 0.08	0.72 ± 0.09	3.59 ± 1.01
020370467	LS	0.186	0.626	50 ± 24	0.78 ± 0.18	1.31 ± 0.25	0.54 ± 0.97
020386743	LS	0.186	0.639	203 ± 8	0.64 ± 0.07	2.68 ± 0.21	1.15 ± 0.42
020461235	LS	0.186	0.597	332 ± 3	0.57 ± 0.02	3.96 ± 0.12	0.47 ± 0.11
020461893	LS	0.186	0.634	283 ± 3	0.46 ± 0.04	2.66 ± 0.11	1.04 ± 0.23
020465775	LS	0.186	0.623	146 ± 3	0.51 ± 0.03	4.04 ± 0.16	0.09 ± 0.17
140083410	12K	0.204	0.958	311 ± 10	0.67 ± 0.09	1.93 ± 0.19	1.90 ± 0.82
140096645	12K	0.204	0.747	217 ± 114	0.93 ± 0.28	1.80 ± 0.40	0.85 ± 1.26
140123568	12K	0.204	0.968	227 ± 229	0.80 ± 1.14	1.39 ± 0.98	0.08 ± 32.82
140137235	12K	0.204	0.920	116 ± 24	0.47 ± 0.31	5.45 ± 2.03	0.88 ± 1.73
140217425	12K	0.204	0.713	260 ± 1	0.25 ± 0.01	8.96 ± 0.13	0.33 ± 0.04
140258511	12K	0.204	0.768	265 ± 42	0.54 ± 0.51	2.62 ± 2.11	0.40 ± 3.22
140262766	12K	0.204	0.613	142 ± 28	0.65 ± 0.26	2.07 ± 0.70	0.63 ± 1.34
140545062	12K	0.204	0.724	216 ± 41	0.46 ± 0.35	2.89 ± 6.86	0.06 ± 4.03
220014252	LS	0.186	0.687	136 ± 2	0.22 ± 0.06	3.32 ± 0.12	0.94 ± 0.23
220015726	LS	0.186	0.686	193 ± 12	0.78 ± 0.07	2.76 ± 0.25	0.26 ± 0.35
220071601	LS	0.186	0.665	74 ± 1	0.31 ± 0.01	8.13 ± 0.13	0.21 ± 0.04
220148046	LS	0.186	0.800	236 ± 10	0.10 ± 0.87	1.98 ± 0.38	2.10 ± 2.36
220376206	LS	0.186	0.749	235 ± 1	0.46 ± 0.01	5.29 ± 0.08	0.21 ± 0.05
220386469	LS	0.186	0.725	168 ± 7	0.70 ± 0.05	2.90 ± 0.13	0.73 ± 0.26
220397579	LS	0.186	0.707	331 ± 5	0.43 ± 0.06	3.02 ± 0.32	0.21 ± 0.37
220397579s	LS	0.186	0.707	332 ± 2	0.14 ± 0.03	9.77 ± 3.29	3.77 ± 1.49
220544103	LS	0.186	0.686	208 ± 55	0.08 ± 0.16	5.61 ± 19.89	0.02 ± 0.91
220544103s	LS	0.186	0.686	161 ± 5	0.26 ± 0.06	5.47 ± 0.44	1.67 ± 0.77
220544394	LS	0.186	0.624	234 ± 7	0.70 ± 0.04	3.40 ± 0.15	0.48 ± 0.13
220544394s	LS	0.186	0.624	233 ± 10	0.30 ± 0.17	3.69 ± 0.70	1.18 ± 1.02
220576226	LS	0.186	0.675	237 ± 15	0.85 ± 0.06	2.16 ± 0.12	1.58 ± 0.36
220578040	LS	0.186	0.658	112 ± 14	0.89 ± 0.04	3.81 ± 0.12	0.42 ± 0.10
220584167	LS	0.186	0.684	193 ± 2	0.52 ± 0.01	7.17 ± 0.19	0.14 ± 0.06
220596913	LS	0.186	0.623	255 ± 1	0.18 ± 0.02	9.49 ± 11.54	0.04 ± 0.12
910193711	LS	0.186	0.597	9 ± 3	0.50 ± 0.03	2.27 ± 0.06	0.50 ± 0.17
910279515	LS	0.186	0.602	131 ± 3	0.61 ± 0.02	4.44 ± 0.11	0.28 ± 0.08

Notes. In Col. (2), “LS” refers to the CFHT Legacy Survey (<http://www.cfht.hawaii.edu/Science/CFHLS/>) whereas “12K” refers to the CFH12K/CFHT survey (McCracken et al. 2003). The pixel scale of the CFHT image and the seeing determined from close stars are respectively given in Cols. (3) and (4). The parameters of the model are the position angle of the major axis (5), the axis ratio (6), the effective radius (7), the Sersic index (7) and the magnitude (not given here due to a zero point magnitude mismatch between the two CFHT surveys). The suffix “s” refers to secondary objects which have been detected both in H α and in the I-band image. The companion’s parameters are only given when they are resolved.

the inner parts. Details of the method are given in Appendix A Appendix B and the resulting parameters are given in Table 3. of Epinat et al. (2010). The maps of the models are shown in For non-isolated galaxies (see Sect. 4), we fitted the various

B. Epinat et al.: MASSIV: Mass Assembly Survey with SINFONI in VVDS. II.

Table 3. Physical parameters from kinematic modeling.

VVDS ID	i	PA_k	r_t	V_t	V_t/r_t	Res_V	χ^2	V_{max}	σ	R_{last}
(1)	[°] (2)	[°] (3)	[kpc] (4)	[km s ⁻¹] (5)	[km s ⁻¹ kpc ⁻¹] (6)	[km s ⁻¹] (7)	(8)	[km s ⁻¹] (9)	[km s ⁻¹] (10)	[kpc] (11)
020106882	52 ± 3	317 ± 1	1.1 ± 0.1	133 ± 3	126	12	1.6	133 ± 25	41 ± 31	5.1
020116027	68 ± 4	207 ± 6	4.3 ± 3.0	27 ± 6	6	10	3.7	27 ± 10	47 ± 15	6.5
020147106	60 ± 24	317 ± 3	1.6 ± 4.6	26 ± 3	16	6	1.9	26 ± 51	81 ± 10	7.8
020149061	60 ± 24	235 ± 3	42.0	976	23	15	1.9	112 ± 216	73 ± 20	4.8
020164388	45 ± 4	97 ± 4	7.2 ± 4.9	79 ± 44	11	8	0.9	79 ± 19	52 ± 19	8.2
020167131	53 ± 4	185 ± 10	1.0 ± 0.1	127 ± 43	123	13	0.9	127 ± 29	26 ± 37	1.8
020182331	61 ± 2	234 ± 4	27.0	647	24	11	0.9	132 ± 26	67 ± 29	5.5
020193070	71 ± 4	184 ± 2	108.7	3217	70	15	1.8	117 ± 23	33 ± 27	3.9
020208482	46 ± 3	336 ± 10	1.0 ± 0.1	158 ± 38	157	11	0.6	158 ± 31	7 ± 11	1.4
020214655	66 ± 11	341 ± 4	1.0 ± 0.1	52 ± 4	51	8	1.1	52 ± 14	63 ± 25	5.7
020239133	67 ± 4	109 ± 6	24.6	766	31	13	0.9	149 ± 33	75 ± 35	4.8
020240675	60 ± 24	183 ± 12	8.9	116	13	13	1.5	50 ± 97	33 ± 17	3.8
020255799	37 ± 15	89 ± 76	1.0 ± 0.1	14 ± 16	14	10	0.6	14 ± 26	76 ± 23	4.0
020261328	59 ± 11	179 ± 2	6.2	154	25	10	1.6	127 ± 35	54 ± 19	5.1
020278667	42 ± 20	175 ± 6	3.7	224	60	16	1.5	77 ± 189	52 ± 37	1.3
020283083	68 ± 2	359 ± 4	1.0 ± 0.1	59 ± 5	57	11	1.4	59 ± 12	39 ± 23	5.6
020283830	70 ± 1	156 ± 2	1.9 ± 3.4	186 ± 5	96	21	4.7	186 ± 30	17 ± 24	7.9
020294045	52 ± 5	3 ± 1	31.1	1332	43	33	12.4	234 ± 51	60 ± 45	5.5
020363717	60 ± 24	106 ± 6	16.4	123	8	9	1.3	45 ± 86	91 ± 14	6.0
020370467	39 ± 14	19 ± 11	1.0 ± 0.1	51 ± 10	49	18	1.2	51 ± 64	86 ± 35	5.5
020386743	50 ± 4	136 ± 3	37.0	286	8	9	2.5	42 ± 10	53 ± 18	5.4
020461235	55 ± 1	351 ± 2	1.2 ± 7.7	82 ± 4	67	11	1.8	82 ± 16	24 ± 22	5.4
020461893	63 ± 4	279 ± 4	1.2 ± 9.1	58 ± 6	46	8	1.1	58 ± 13	67 ± 22	6.5
020465775	59 ± 2	178 ± 5	1.1 ± 0.1	68 ± 7	65	13	1.5	68 ± 15	84 ± 30	4.9
140083410	48 ± 16	39 ± 6	1.0 ± 0.1	30 ± 4	30	10	2.3	30 ± 33	66 ± 21	5.3
140096645	22 ± 15	197 ± 1	1.1 ± 1.1	295 ± 12	258	12	6.1	295 ± 709	77 ± 27	4.5
140123568	60 ± 24	184 ± 14	1.0	108	108	6	0.7	50 ± 99	73 ± 24	0.5
140137235	62 ± 1	123 ± 16	2.2	200	91	4	0.6	62 ± 11	17 ± 27	0.7
140217425	76 ± 2	258 ± 1	16.1	499	31	29	20.7	320 ± 46 ¹	45 ± 33	14.5
140258511	57 ± 3	213 ± 1	1.0 ± 0.1	124 ± 4	119	19	5.2	124 ± 26	25 ± 29	5.2
140262766	60 ± 24	175 ± 2	18.1	524	29	9	1.2	119 ± 231	39 ± 16	4.1
140545062	63 ± 5	229 ± 1	13.9	378	27	13	3.6	204 ± 46	67 ± 29	7.5
220014252	77 ± 2	141 ± 1	1.0 ± 0.1	129 ± 1	123	15	4.5	129 ± 27	90 ± 28	10.3
220015726	39 ± 15	186 ± 1	1.5 ± 0.4	231 ± 4	155	10	3.1	231 ± 356	62 ± 22	3.7
220148046	60 ± 24	261 ± 15	1.3	61	49	10	2.9	42 ± 83	46 ± 21	0.9
220376206	63 ± 2	225 ± 1	8.3 ± 0.2	201 ± 4	24	15	4.8	201 ± 27	73 ± 26	10.0
220386469	46 ± 7	151 ± 10	63.0	973	15	16	2.0	40 ± 11	43 ± 25	2.6
220397579	65 ± 6	1 ± 9	35.9	32	1	9	7.0	9 ± 10	59 ± 17	10.2
220397579s	80 ± 5	344 ± 2	3.8 ± 1.2	222 ± 9	59	17	2.6	222 ± 15	27 ± 37	6.2
220544103	80 ± 4	198 ± 1	1.1 ± 0.1	137 ± 2	130	12	4.4	137 ± 24	71 ± 19	7.6
220544394	46 ± 2	180 ± 2	2.0 ± 1.5	55 ± 4	27	5	0.8	55 ± 11	49 ± 17	5.0
220544394s	73 ± 5	198 ± 4	1.0 ± 0.1	94 ± 8	94	16	1.7	94 ± 18	32 ± 28	4.5
220576226	32 ± 6	283 ± 3	1.0 ± 0.1	30 ± 1	30	4	1.1	30 ± 12	51 ± 14	6.1
220578040	27 ± 9	103 ± 2	18.5	654	35	16	5.1	247 ± 205	50 ± 23	7.0
220584167	59 ± 1	178 ± 1	9.7 ± 0.2	234 ± 2	24	15	7.9	234 ± 35	49 ± 21	13.1
220596913	80 ± 2	247 ± 1	0.9 ± 0.1	141 ± 2	153	22	4.3	141 ± 10	38 ± 28	9.3
910193711	60 ± 8	39 ± 4	3.5 ± 0.7	63 ± 8	18	17	2.6	63 ± 12	80 ± 37	4.1
910279515	52 ± 4	108 ± 7	3.8	265	70	34	5.3	186 ± 14	47 ± 40	2.7

Notes. The parameters of the kinematics modeling are the inclination (2), the position angle of the major axis (3), the turnover radius (4) and velocity (5). The inner slope is given in Col. (6). The mean residual of the velocity field is given in Col. (7) and the fit chi square in Col. (8). The maximum rotational velocity (9) and the mean velocity dispersion corrected from beam smearing (10) are computed after from the results of the fit. The extent of the velocity field (11) is derived using a S/N threshold of 3. The suffix “s” refers to secondary objects which have been detected in $H\alpha$ and large enough to perform kinematics modeling. No error is given for r_t and V_t when the plateau is not reached (see Sect. 3.2.3). ⁽¹⁾ V_{max} is not coming from the model but from a detailed analysis (see Appendix A).

components separately when they were sufficiently extended (VVDS220397579). In the table, companions have the suffix “s”.

The model rotation curve only reproduces the velocity field and is not based on a gravitational potential model (as for instance in Förster Schreiber et al. 2006; Cresci et al. 2009; Gnerucci et al. 2011) since the shape of the gravitational potential is unknown and can probably not be described by a stellar component only. Note that a gaseous thin disk in rotation is not

incompatible with a spheroidal stellar distribution such as was observed in local ellipticals (e.g. Sarzi et al. 2006) or as suggested from numerical simulations (e.g. Bournaud et al. 2007; Xu et al. 2010).

As described in the following sections, some parameters are difficult to constrain from the kinematics. To reduce the number of free parameters, we constrained the center and the inclination from the morphology to model the kinematics, assuming that

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the stars and the ionized gas follow a common distribution. Thus only four parameters remain free and can be reasonably constrained from our observed velocity fields. Using these constraints, [Epinat et al. \(2010\)](#) have shown from 137 galaxies of the GHASP sample (the largest 2D kinematics sample of nearby late-type galaxies described in [Epinat et al. 2008b,a](#)) projected at $z \sim 1.7$, that this method enables one to recover the other model parameters statistically.

3.2.1. Center

Owing to the low spatial resolution of our data, the kinematic center of a given galaxy is barely constrained from the kinematics ([Epinat et al. 2010](#)). Therefore the centers are taken from the *I*-band morphology since we expect that in the inner parts of a galaxy at these redshifts, the stars dominate the gravitational potential.

Thanks to the method we used, the astrometry in the SINFONI data cubes matches the *I*-band astrometry. However, a post correction was applied to match the outer isophotes of the galaxy in the CFHT images and in the $H\alpha$ maps. The median offset of this post correction is $0.18''$, which agrees with the accuracy of the SINFONI pointing system, which is estimated to be $0.1\text{--}0.2''$ (see SINFONI manual, using a guiding star instead of offsetting from a bright star). This is also the final accuracy of our astrometry. This offset was computed for 38 galaxies. The other galaxies were excluded (i) when no PSF star was observed or (ii) when they were observed with AO because the observing sequence did not allow us to compute an astrometry correction. We find a good agreement between the centers derived from the *I*-band images and the peak in the $H\alpha$ maps.

3.2.2. Inclination

Disk inclination is a critical parameter to estimate because it is directly linked to the rotational velocity (V_θ) of the disks. Indeed, we measured the velocities projected along the line of sight: $V_{\text{los}} = V_\theta \times \sin i$. Consequently, the observed velocity has to be corrected for the inclination and this correction is larger for galaxies with low inclination. Owing to this degeneracy between inclination and rotation velocity in rotating disk models, the inclination is not well constrained from the kinematics alone. This degeneracy can theoretically be solved for high-resolution observations but not for observations with strong beam-smearing ([Epinat et al. 2010](#)): this would lead to an erroneous rotational velocity. Therefore the inclination was constrained from the morphology axis ratio. In addition, since the uncertainty on the inclination (its sine) directly impacts the uncertainty on the rotation velocity, we paid special attention in deriving realistic error bars on the inclination. The thickness of the disk, which is considered as null in our models, could impact the determination of the inclination, mainly for edge-on galaxies. However, for these objects, the deprojection has a negligible impact on the determination of V_{max} (less than 3% considering a thickness leading to $b/a \sim 0.2$).

To take into account both the uncertainty on the model and the uncertainty on the PSF used to recover galaxy parameters in GALFIT, we used a Monte Carlo method. For each galaxy we simulated 2000 synthetic sources with the same structural parameters (magnitude, Sersic index, effective radius, position angle), but a random value of b/a . The PSF to simulate these sources was also randomly chosen among different stars in the fields. Poisson noise was added and the simulated sources were

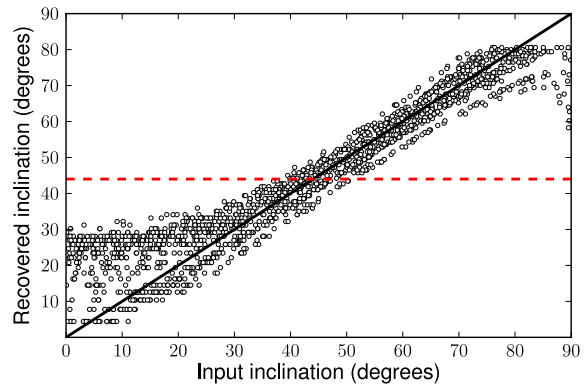


Fig. 1. Example of the method used to estimate the uncertainty on the inclination for galaxy VVDS020164388. The *x*-axis corresponds to the input inclination of the modeled galaxies, the *y*-axis corresponds to the inclination recovered by GALFIT for each modeled galaxy and the red dashed line corresponds to the inclination of the real galaxy. The uncertainty is measured along this line as the range that contains 68% of the points (1σ).

placed in a nearby piece of sky from the original image (three pieces for each galaxy). The simulated sources were then modeled with GALFIT using the PSF used for the real galaxy. The final uncertainty was estimated from the range of input inclinations that led to the inclination that was measured with GALFIT on the real galaxy (see Fig. 1). For the smallest galaxies, the inclination was not constrained enough. We therefore decided to use an inclination of 60° , the median value for randomly distributed disks, and an uncertainty of 24° , which yields a probability of 0.68 (1σ).

We also checked that there was no bias by studying the inclination distribution for the MASSIV sample. Theoretically, for uniformly randomly oriented thin disks (i.e. with a null thickness) the probability to observe a disk with an inclination between θ_1 and θ_2 is equal to $|\cos \theta_1 - \cos \theta_2|$, leading to distribution as displayed in black in Fig. 2 with a median value of 60° . In Fig. 2, we observe that the distribution of MASSIV galaxies with constrained inclinations (filled blue histogram) misses face-on objects ($i = 0^\circ$) and edge-on objects ($i = 90^\circ$) and has an excess of objects with intermediate inclinations compared to the theoretical distribution. The most plausible explanation for the lack of edge-on morphologies is that disks are thick. It could also be that extinction is higher in these galaxies, inducing a non-detection of the $[\text{O II}] \lambda 3727$ line in the VVDS spectra. Figure 1 clearly shows that GALFIT cannot recover extreme inclinations. However, the lack of face-on objects could also be attributed to the small numbers expected or to a computing bias. It could also be that galaxies do not have regular morphologies and that, due to surface brightness dimming, one can only observe the clumpy irregular emission. Indeed, *I*-band morphology could be in some cases contaminated by gaseous emission lines like $[\text{O II}] \lambda 3727$ because the MegaCam/MegaPrime *i'* band filter covers the 0.70 to $0.84 \mu\text{m}$ spectral range, which corresponds to $[\text{O II}] \lambda 3727$ redshifted between 0.87 and 1.26. Unfortunately, the CFHT *I*-band imaging resolution does not allow us to conclude about the clumpy emission. The peak around 60° could also be attributed to (i) the thickness of disks and (ii) the observation of galaxies that are not disks (mergers for example). The median of the distribution for our sample (59°) is compatible with the median expected for a distribution of disks with random inclination.

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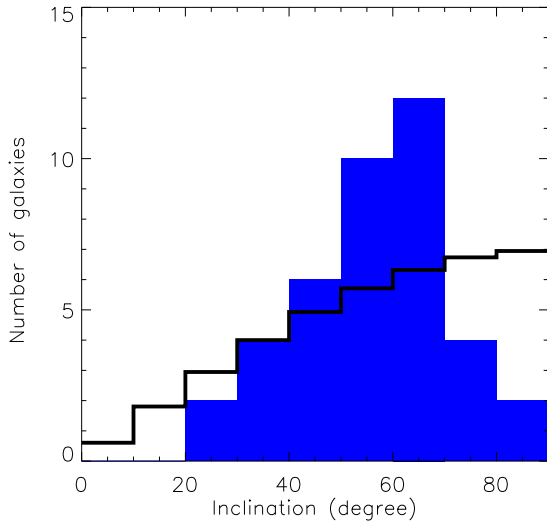


Fig. 2. Distribution of the inclinations for the 40 galaxies of the MASSIV sample (with a constrained inclination) using a Sersic profile with GALFIT and assuming they are thin disks (filled blue histogram) compared with the theoretical distribution for a randomly oriented thin disk (black histogram).

3.2.3. Constraints on the other model parameters

Since inner velocity gradients reaching a plateau (V_t) within less than one pixel could not be resolved with our spatial resolution which is higher than four pixels, the turnover radius r_t was constrained to be at least one pixel during the fitting process. However, when the model converged toward this value, the statistical error becomes null. On the other hand, when the plateau is not reached according to the χ^2 minimization, the error on r_t becomes large and so does the error on V_t because the models mainly constrain the slope. Therefore the errors on these parameters are difficult to interpret and to propagate to compute the error on the inner slope of the velocity gradient. A solution would consist in using a model described by the slope α and the turnover radius r_t to estimate the error on the slope. This test has been performed but did not give more realistic errors in many cases. Because this solution was not convincing, we used the original model. For these galaxies, no error is indicated for r_t and V_t in Table 3. Although r_t is not well constrained for small galaxies, Epinat et al. (2010) showed with 137 galaxies projected at $z \sim 1.7$ that leaving r_t as a free parameter statistically gives a good estimate of the shape of the rotation curve. It is therefore necessary to let r_t be a free parameter, in particular for the largest galaxies.

3.2.4. Parameters deduced from the kinematic maps and models

Radius of the H α extent R_{last}

This radius is computed from the cleaned maps. The center and position angle from the kinematic best-fit models were used to derive the radius of each pixel. R_{last} is the radius that both sides of the galaxy reach.

Velocity shear V_{shear}

V_{shear} was computed as the total shear observed in the modeled velocity field shown in Appendix B (i.e. with an S/N threshold of 3) which takes into account the uncertainty on

the observed velocity field. It is a projected velocity along the line of sight and is not inclination-corrected and consequently makes no assumption on the geometry of the galaxy. The use of the model enables us to smooth the velocity field at the outskirts.

Maximum rotation velocity V_{max}

This was computed according to the model at the R_{last} radius. Two sources of uncertainties were added in quadrature to compute the final uncertainty on V_{max} .

The first one is the uncertainty on the inclination (cf. Sect. 3.2.2). To propagate uncertainties from inclination to rotation velocity, we used a Monte Carlo method assuming a Gaussian distribution for the inclination. The uncertainty on V_{max} was the standard deviation of the resulting distribution on this parameter.

The second source of uncertainty is related to the modeling. Because the uncertainty on V_t is a statistical one, it can be fairly small. Instead of using an approach based on these errors, the GHASP sample (Epinat et al. 2008a,b) was exploited to compute model uncertainties related to the size of the galaxies with respect to the seeing. Using the same method as Epinat et al. (2010), 136 GHASP galaxies were projected at $z = 1.33$ under typical seeing and sampling of SINFONI observations of the MASSIV sample and the S/N of the simulations was adjusted to match the H α fluxes from the MASSIV sample. Figure 13 of Epinat et al. (2010) was reproduced with these new simulations. This figure displays the relative error on the maximum velocity determination with respect to the beam-smearing parameter defined as $B = D_{25}/2s$, where D_{25} is the optical diameter and s is the seeing FWHM. A linear regression was performed to estimate the evolution of the accuracy of the fit with galaxy size. The best fit gives

$$\delta V_{max} = V_{max} \times \frac{27.5 - 5.8B}{100}. \quad (2)$$

The percentage of galaxies below this relation is about 60%. Assuming a Gaussian distribution, $1-\sigma$ corresponds to a confidence level of 68%. We therefore simply approximated that δV_{max} is the $1-\sigma$ uncertainty. For the MASSIV sample, B was estimated from half-light radii (R_e) determined by GALFIT on I -band images (see Sect. 3.1). Indeed, assuming an exponential disk, the following relation can be written: $D_{25}/2 = 1.9R_e$. The model uncertainty was consequently determined following Eq. (2). A minimum uncertainty of 10 km s^{-1} was imposed, however. This approach is well-adapted for rotators because it was built from a control sample of rotators.

Local velocity dispersion σ

The velocity dispersion was computed as in Epinat et al. (2009): this is the average of the velocity dispersion map corrected for beam-smearing effects (see Sect. 3.2) and spectral PSF. A weight proportional to the inverse of the uncertainty on the velocity dispersion (estimated before correcting for beam-smearing and spectral resolution) was attributed to each pixel to compute the average. The uncertainty on this parameter was computed as the weighted standard deviation.

4. Kinematics classification

To distinguish between galaxy formation scenarios, it is necessary to know the dynamical state of galaxies for large samples and at various redshifts.

4.1. Previous classification schemes of large samples

The first kinematics classification of distant galaxies observed using integral field unit techniques was provided by Flores et al. (2006) in the frame of the IMAGES sample ($0.4 < z < 0.75$) which contains 68 classified galaxies (Neichel et al. 2008). Their classification is a visual one that relies on both HST optical images and FLAMES/GIRAFFE data. The galaxies were split into three classes: (i) rotating disks whose morphological and kinematic position angles match well and with a velocity dispersion peak in the center, (ii) perturbed rotators whose morphological and kinematic position angles match well but that have a peak in the velocity dispersion offset from the center and (iii) complex kinematics systems with both a disagreement between morphological and kinematic position angles and a peak in the velocity dispersion offset from the center. The IMAGES survey has a majority of galaxies with complex kinematics (44%), and a nearly equivalent fraction of rotating disks (29.5%) and of perturbed rotators (26.5%) (Yang et al. 2008; Neichel et al. 2008). From these numbers, they concluded that at that epoch merging is still more active than in the local Universe.

The classification of the SINS sample ($z \sim 2.2$) relies on the asymmetries measured on both the velocity and velocity dispersion fields to distinguish rotator-like from merger-like galaxies. These asymmetries were derived either from a kinemetry analysis (Shapiro et al. 2008) or from a qualitative assessment (Förster Schreiber et al. 2009). Then, comparing the mean local velocity dispersion and the rotational velocity, these galaxies (both mergers and rotators) were classified either as rotation-dominated ($V_{\max}/\sigma_0 > 1$) or dispersion-dominated ($V_{\max}/\sigma_0 < 1$). For galaxies with small size or low S/N , they instead compared the full velocity shear v_{obs} with the integrated line width σ_{int} and used a threshold of $v_{\text{obs}}/(2\sigma_{\text{int}}) \sim 0.4$. On the one hand, Förster Schreiber et al. (2009) found that one third of the 62 galaxies in the H α SINS sample contains mergers. On the other hand, they also found that one third of this sample contains rotation-dominated systems, another third corresponds to dispersion-dominated systems, and the last third are not classified. The authors interpret the large amount of dispersion-dominated disks as an evidence for cold gas accretion along cosmic web filaments.

For the LSD/AMAZE sample, the classification only distinguishes rotator-like from perturbed galaxies because galaxies have a fairly low S/N . This classification relies on the modeling of the velocity field as an inclined X - Y plane (Gnerucci et al. 2011). The criterion is based on the χ^2 of the fit by the plane and on a constraint on the reliability of the inclination of the plane.

In the pilot run study of MASSIV (Epinat et al. 2009), which was based on a visual kinematic classification, nine galaxies with $1.2 < z < 1.6$ were classified into three groups: (i) three galaxies are mergers, showing disturbed kinematics and possibly several components, (ii) two objects are classified as rotation-dominated disks, with clear signs of rotation and with a maximum rotation velocity higher than the mean local velocity dispersion and (iii) four galaxies are dispersion-dominated disks, with clear signs of rotation but with a maximum rotation velocity lower than the mean local velocity dispersion. It was concluded that about one third of these galaxies are observed during some merging event, whereas one third are rotation-dominated and the other third are dispersion-dominated.

In this paper, we present a new kinematic classification for the MASSIV sample to put constraints on the dynamical state of galaxies at $z \sim 1.2$. The classification scheme was developed in several steps. First, eight people of the collaboration independently defined their own criteria. These criteria were

then reconciled and discussed during a common session and led to a unique classification based both on the close environment and on the velocity shear strength. The visual classification helped in defining measurable criteria to build an automatic, thus reproducible, classification (this led to minor changes in the final classes). The final classification was also refined to describe the dynamical state and support of the galaxies.

4.2. Galaxy small-scale environment

The study of galaxy close environment is crucial for inferring a merger rate on the population probed by the MASSIV sample at $z \sim 1.2$. Using SINFONI data and CFHT imaging, galaxies could be classified as interacting or isolated. Flags have also been attributed to this classification to qualify its reliability. Flags “A”, “B” and “C” mean secure classification (>90% probability), confident (~75% probability) classification and poor reliability (~50% probability), respectively.

The SINFONI field-of-view was explored to detect emission lines that might be attributed to companions. The same field was also explored systematically in the I -band images. Owing to the nodding strategy, the field-of-view is larger for seeing-limited observations. In that case, the shape of the field results from the superposition of two 8" square fields overlapped on their opposite corners (4" quadrants). Thus, the field-of-view of the combined data around each target can be as large as 6" (corresponding to ~50 kpc at $z \sim 1.2$). In the non-overlapping regions, the exposure time is half the total on-source exposure time but is sufficient to at least detect emission lines of objects. The explored field is not a full 12" \times 12" square field but only 78% of this area. For AO observations, the field of view is restricted to 3" (corresponding to ~25 kpc), which is insufficient for studying the environment and therefore only the I -band image was explored to see if a possible companion was observed in the 6" around the main object.

Galaxies were classified as interacting with a confidence flag “A” when the following conditions were met: (i) the average S/N (see Table 4) has to be higher than three in an area larger than the seeing, (ii) the systemic velocity difference between the two components has to be lower than 1000 km s⁻¹ and (iii) an optical counterpart has to be observed within the CFHT I -band image. If the companion was extended but slightly smaller than the seeing, we assigned a flag “B”. When no counterpart was observed in I -band or if some galaxies were observed in the neighborhood in I -band but not in the SINFONI datacube, the object was considered as isolated with a flag “B”. A flag “C” was assigned for AO observations unless absolutely no other galaxy was observed in I -band around the object. This is detailed in Appendix A.

In some cases, the morphology is elongated and several H α blobs can be detected along the elongation, sometimes also in I -band images (see comments in Appendix A). It is not straightforward to assess if these blobs are small objects that are accreted and not star forming regions within one unique galaxy. A monotonic velocity field along the direction of the major axis is likely evidence for rotation. However, if the elongation is the result of the projection of the two components and if the two objects are small (no gradient detectable in each component), this could mimic a monotonic velocity field caused by the difference of radial velocity between the components. Some additional features can also support a merger hypothesis: (i) the angular separation is larger than 3" (i.e. ~25 kpc), because galaxies this large are not expected at high redshift, nor are they observed in the local Universe; (ii) the radial velocity gap is larger than ~600 km s⁻¹, because this rotational velocity would imply an

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unrealistic dynamical mass for a unique object. If one of these features was seen, the galaxies were classified as interacting but with a flag “B”. If none of them was observed, the galaxies were classified as isolated but a flag “B” was assigned. For the difficult cases, we were able to use the morphology traced by old stars (mainly observable in near-infrared bands): the old star population may not follow the star forming regions in case of unique objects. The S/N of these J - or H -band continuum morphologies recovered from SINFONI data is fairly low, however. It is therefore not straightforward to arrive at definite conclusions.

When components are very close, i.e. about to merge, the situation is more complex. However, perturbed line profiles (e.g. VVDS020294045, cf. Appendix A) in the region between the two blobs favor a strong discontinuity in the velocity field, which points toward a two-component hypothesis. These cases received a flag “B”.

For ongoing mergers one expects to observe peculiar kinematic signatures. Therefore, we classified some galaxies as interacting because their velocity fields, velocity dispersion maps and morphologies were presenting perturbations in the same positions, suggesting for instance the presence of tidal arms (e.g. VVDS020283083, cf. Appendix A). Since these signatures are more subjective, these galaxies usually have a flag “B” for the isolation criterion.

Other galaxies were classified as isolated. However, when there were some perturbations in the velocity field or a peak in the velocity dispersion map, this could be interpreted as signs for merger remnants, but because it concerns the galaxy outskirts, it is more likely noise and we assigned a flag “B”.

4.3. Velocity shear strength

For the kinematic classification, we used a first criterion as simple and objective as possible based on the velocity shear V_{shear} (see Sect. 3.2.4). The sample was accordingly divided into low-velocity shear galaxies ($V_{\text{shear}} < 100 \text{ km s}^{-1}$) and high-velocity shear galaxies ($V_{\text{shear}} > 100 \text{ km s}^{-1}$). This gives a rough idea about the dynamical state of a galaxy without taking into account more complex motions. For VVDS220397579 and VVDS220544394 it was possible to study the velocity shear of the companion as well.

This criterion has the advantage that it can be easily measured. However, the value of V_{shear} for a given galaxy may vary with the orientation. This is particularly true for galaxies in ordered rotation (V_{shear} is lower when observed face-on). However, the distribution of the inclinations, as seen in Fig. 2, shows that only eight galaxies according to the theoretical distribution and only three according to the observed distribution may need a correction larger than a factor of two (corresponding to an inclination of 30°) for the measurement of the velocity shear.

In galaxies with a low-velocity shear there might be several classes of objects: (i) very low-mass objects; (ii) face-on rotating galaxies; (iii) ongoing mergers in a transient state; (iv) spheroids, if one expects that the gas in these objects follows the distribution of stars; and (v) galaxies with a non/slowly rotating gaseous component.

4.4. Rotating and non-rotating galaxies

For a specific analysis (e.g. Tully-Fisher relation, see Vergani et al. 2012), it is essential to be able to define a robust sample of rotating galaxies. We expect a rough agreement of the morphological and kinematic position angles for rotators. They

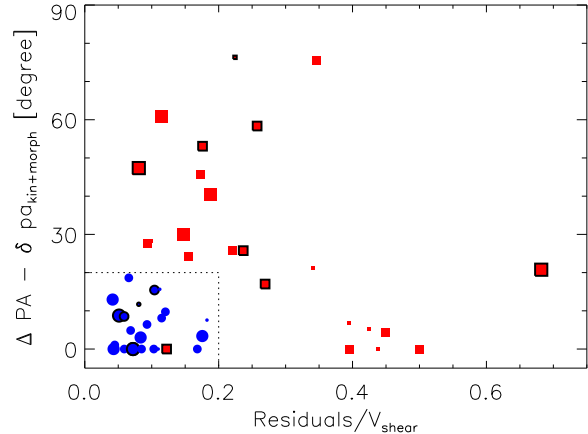


Fig. 3. Disagreement between morphological and kinematic position angles (within the errors) as a function of the velocity field residuals normalized by the velocity shear. Blue dots and red squares represent rotating and non-rotating galaxies. Symbols with black contours mark interacting galaxies. The symbol size is related to the S/N of the SINFONI data: small, medium and large symbols refer to galaxies with $S/N < 5$, $5 < S/N < 10$ and $S/N > 10$, respectively. One galaxy is not in the expected region (VVDS020294045, cf. Appendix A).

can be slightly different due to perturbations like bars or strong spiral arms. We also expect for these galaxies that the rotation motions dominate over perturbations. To distinguish rotators from non-rotating galaxies, we made a diagram in Fig. 3 with two quantities that quantify these two arguments: the disagreement between morphological and kinematic position angles derived from the models presented in Sects. 3.1 and 3.2 within the uncertainties, $|PA_m - PA_k| - \sqrt{\delta PA_m^2 + \delta PA_k^2}$, and the mean weighted velocity field residuals normalized by the velocity shear. Rotators are therefore defined as galaxies with a position angle difference lower than 20° and velocity field perturbations lower than 20%. These thresholds were chosen from the distribution seen in Fig. 3 because they enable us to isolate a clear cloud of points near the origin (the rotators). We note that these values are also representative of relatively small deviations from rotational motion and position angles. There is one exception in this diagram: VVDS020294045 was classified visually as non-rotating although it lies in the rotator’s region of the diagram. Indeed, the velocity shear visible in its velocity field is probably caused by a very close companion that thus mimics a rotating disk velocity field. The shape of the line suggests a merger. If we were to exclude the companion, which is overlapping with the main galaxy, the resulting velocity field would probably be classified as non-rotating (see Appendix A).

We checked the agreement between this classification and the classification based on kinemetry used in the SINS sample (Shapiro et al. 2008). The agreement is reasonable but we find that galaxies that would be classified as mergers according to their criteria are mainly those that we have classified as non-rotating. Indeed, the classification of Shapiro et al. (2008) relies on the position in a diagram in which galaxies are placed according to their asymmetric velocity and velocity dispersion components. Galaxies show a strong correlation between these two quantities. This correlation arises because asymmetries are normalized by the global velocity shear. We are able to show that the normalized residuals of the model are well-correlated

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with the normalized asymmetries in the velocity field. This is true for large and small galaxies. Thus, our classification adds the agreement between the position angles in morphology and kinematics as a new criterion. One advantage of this criterion is that it is not restricted to galaxies observed with a good S/N and with many resolution elements.

4.5. Dynamical support

The last criterion is related to the dynamical support of galaxies. Indeed, it has been shown that at high redshift ($z > 2$), galaxies are on average more dispersion-dominated than in the intermediate ($z \sim 0.5$) and local Universe. One of our goals is to determine the evolution of the fraction of dispersion-dominated systems with redshift since MASSIV might probe a transitional redshift range.

To quantify the dynamical support, we assumed that all galaxies have some rotation and that this rotation is adequately described by the rotating-disk model. Then, we additionally assumed that the local velocity dispersion of the gas is representative of the random motions in the galaxy. Thus, the dynamical support is estimated by the ratio of the maximum rotation velocity over the local velocity dispersion V_{\max}/σ . This description is intrinsically better justified for rotators, however.

The various classes and related parameters are summarized in Table 4.

5. Discussion

Using the classification scheme described above, we were able to classify as rotating or non-rotating the 46 galaxies (including one companion) in the redshift range $0.9 < z < 1.6$ for which some emission line was detected in the SINFONI datacubes. For part of the discussion below we limit the sample to the 36 galaxies with an average S/N (defined in Sect. 2.4 and reported in Table 4) higher than 5 for which the kinematic classification is more robust (we note that these low S/N galaxies are the majority of galaxies with $R_{\text{last}}/\text{seeing} < 1$, which means that the kinematic model might be less robust). Accordingly among an initial sample of 46 secure galaxies with $0.9 < z < 1.6$, 22% are not robustly classified or not classified from their dynamical properties. This proportion is on the same order for the LSD/AMAZE (Gnerucci et al. 2011). In the MASSIV sample, these galaxies have on average lower stellar masses and star formation rates (see Table 5), the latter explaining the lower S/N of these observations. We also emphasize that the high dispersions are caused by the two galaxies observed with AO, which have larger stellar masses and SFR than the rest of the unclassified galaxies. For those two galaxies the low S/N is due to the small pixel size rather than an intrinsically low SFR. Concerning the close environment criterion, the 41 systems classified with flags “A” and “B” are considered.

5.1. Rotating disks vs. non rotating galaxies

Rotating disks represent at least 44% (20 out of 46 detected galaxies) of the MASSIV “first epoch” sample and about 56% of the high S/N sample. We find a lower percentage of secure non-rotating systems (35% or about 44% of the high S/N sample) that have no observed rotation in the gaseous component or have very disturbed kinematics with respect to their broad band morphology. The physical properties of these two classes of galaxies are on average different. We have performed

Kolmogorov-Smirnov tests (Press et al. 1992) on stellar mass, SFR, half-light radius and velocity dispersion distributions. The two populations look different in terms of stellar mass and SFR at a level slightly higher than 1σ (the probability to follow the same distribution is lower than 0.2). Distributions for size and velocity dispersion are fully compatible. Rotating objects are on average more massive ($M_{\text{star}} = 4.0 \times 10^{10} M_{\odot}$), more star-forming ($SFR = 60 M_{\odot} \text{ yr}^{-1}$) and have larger radii ($R_e = 3.8 \text{ kpc}$) than non-rotating ones ($M_{\text{star}} = 1.6 \times 10^{10} M_{\odot}$, $SFR = 39 M_{\odot} \text{ yr}^{-1}$ and $R_e = 2.7 \text{ kpc}$). On the other hand, these two types of galaxies have very similar velocity dispersions around 60 km s^{-1} (see Table 5). At $z \sim 2.2$, Förster Schreiber et al. (2009) already noticed a similar trend from the SINS sample: dispersion-dominated systems are on average smaller than rotation-dominated ones.

The percentage of rotating systems is higher in MASSIV at $z \sim 1.2$ than in the LSD/AMAZE sample at $z \sim 3.3$. Indeed, Gnerucci et al. (2011) found a lower limit of 34% of rotators in their sample (11 out of 32 detected galaxies that they were able to classify). Their classification is closest to ours since it relies both on velocity field modeling and on the agreement between morphological and kinematic position angle of the major axis. However, owing to their small statistics and their different selection function it cannot be excluded that these proportions are compatible. At $z \sim 2.2$, the comparison with the SINS sample is more difficult because of the selection function, as for LSD/AMAZE, but also because of the classification scheme, which mainly relies on a kinemetry analysis and on a visual inspection (Förster Schreiber et al. 2009). Förster Schreiber et al. were able to unambiguously identify 18 rotating systems (discussed in Cresci et al. 2009), i.e. a lower limit of 35% (18 out of 52 detected systems). These authors were also able to identify 14 dispersion-dominated systems (27% of the SINS sample) based on V_{\max}/σ ratio. Considering only the reliably classified systems, SINS and MASSIV percentages are similar. At lower redshift, the results of the IMAGES sample ($z \sim 0.6$) show that 63% of the sample show signs of rotation (both rotating disks and perturbed rotators classes, Puech et al. 2008). This percentage is higher than for MASSIV and favors an interpretation in which gas in star-forming systems is stabilizing into disks while the Universe evolves. This conclusion has to be balanced with the various selection functions and classification methods, in the same way as for the comparison with the other surveys discussed previously. Concerning the selection, the various authors claim that they observed representative sets of star-forming galaxies in each redshift range.

In the MASSIV sample, the most massive galaxies ($\log M_{\text{star}} > 10.5$) are mainly identified as disks in rotation. However, this is true for the gaseous phase and does not necessarily imply that the stars are settled into a disk. The best spheroidal candidates may be the most compact and roundest massive galaxies. Indeed, elliptical galaxies can be flattened, too, but should be rounder than disks on average. There are four potential candidates even if only one of them has an axis ratio fully compatible with zero (see Fig. 4). These may also be nearly face-on disks if observed velocity shear is very low. In addition, given our definition of rotators, it is highly probable that gas and stars share a common disk since both morphological and kinematic major axis agree well. The morphology obtained from the CFHT imaging used for MASSIV has an insufficient low spatial resolution to allow addressing this problem unambiguously but it is worth noticing that in the local Universe stellar disks are observed in elliptical galaxies (e.g. Kuntschner et al. 2010).

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Table 4. Kinematics and close environment classification of MASSIV “first epoch” sample galaxies.

VVDS ID (1)	S/N (2)	$\frac{R_{\text{last}}}{\text{Seeing}}$ (3)	ΔPA [°] (4)	$\frac{R_{\text{res}}}{V_{\text{shear}/2}}$ (5)	V_{max}/σ (6)	Shear (7)	Rotator (8)	Isolated (9)	Isolation flag (10)
020106882	7.6	1.2	23 ± 4	0.07	3.2	High	Yes	Yes	B
020116027	8.2	1.3	23 ± 6	0.27	0.6	Low	No	No	A
020126402	<3	–	–	–	–	–	–	–	–
020147106	12.2	1.4	7 ± 3	0.18	0.3	Low	Yes	Yes	B
020149061	6.2	0.7	34 ± 90	0.17	1.5	Low	Yes	Yes	B
020164388	11.0	1.2	65 ± 5	0.11	1.5	Low	No	Yes	B
020167131	3.7	0.3	87 ± 11	0.22	5.0	Low	No	No	A
020182331	6.1	0.9	34 ± 6	0.09	2.0	High	No	Yes	B
020193070	4.9	0.8	31 ± 3	0.10	3.6	High	No	Yes	B
020208482	4.1	0.3	20 ± 12	0.18	22.9	Low	Yes	Yes	B
020214655	7.5	0.8	51 ± 5	0.17	0.8	Low	No	Yes	B
020217890	<3	–	–	–	–	–	–	–	–
020239133	5.3	0.8	14 ± 6	0.11	2.0	High	Yes	Yes	B
020240675	5.5	0.5	7 ± 18	0.40	1.5	Low	No	Yes	B
020255799	4.5	0.6	75 ± 77	1.04	0.2	Low	No	Yes	B
020261328	7.8	1.0	8 ± 8	0.07	2.4	High	Yes	Yes	B
020278667	4.1	0.2	35 ± 14	0.34	1.5	Low	No	Yes	C
020283083	7.0	0.9	58 ± 5	0.18	1.5	Low	No	No	B
020283830	4.8	1.2	14 ± 2	0.08	11.3	High	Yes	No	B
020294045 ¹	7.0	1.2	2 ± 4	0.12	3.9	Low	No	No	B
020306817	<3	–	–	–	–	–	–	–	–
020363717	11.4	1.1	49 ± 8	0.19	0.5	Low	No	Yes	B
020370467	5.5	0.9	31 ± 27	0.45	0.6	Low	No	Yes	B
020386743	8.8	0.9	67 ± 8	0.26	0.8	Low	No	No	A
020461235	5.8	1.1	19 ± 4	0.10	3.5	High	Yes	No	B
020461893	7.0	1.3	4 ± 5	0.10	0.9	Low	Yes	Yes	B
020465775	6.8	0.7	31 ± 5	0.24	0.8	Low	No	No	B
140083410	6.0	1.0	88 ± 12	0.35	0.4	Low	No	Yes	A
140096645	10.3	1.0	20 ± 90	0.07	3.9	High	Yes	No	B
140123568	3.5	0.1	43 ± 90	0.44	0.7	Low	No	Yes	B
140137235	3.2	0.1	8 ± 29	0.11	3.5	Low	Yes	Yes	B
140217425	6.8	1.9	2 ± 1	0.04	7.1	High	Yes	Yes	B
140258511	8.1	1.3	52 ± 42	0.12	5.1	High	Yes	Yes	A
140262766	6.5	1.0	32 ± 28	0.07	3.1	High	Yes	Yes	B
140545062	8.5	1.3	14 ± 41	0.06	3.1	High	Yes	Yes	B
220014252	11.0	1.7	5 ± 2	0.08	1.4	High	Yes	Yes	B
220015726	10.4	1.0	8 ± 12	0.04	3.7	High	Yes	Yes	B
220071601	<3	–	–	–	–	–	–	–	–
220148046	4.0	0.4	25 ± 18	0.39	0.9	Low	No	Yes	C
220376206	12.6	2.4	10 ± 1	0.05	2.8	High	Yes	No	B
220386469	4.5	1.1	17 ± 12	0.42	0.9	Low	No	Yes	B
220397579	15.7	2.0	31 ± 10	0.68	0.2	Low	No	No	A
220397579 _s	5.4	1.2	11 ± 3	0.06	8.3	High	Yes	No	A
220544103	10.4	1.2	11 ± 55	0.07	1.9	High	Yes	No	B
220544394	10.5	1.1	54 ± 7	0.08	1.1	Low	No	No	A
220576226	11.6	1.3	45 ± 15	0.15	0.6	Low	No	Yes	B
220578040	5.9	1.4	10 ± 14	0.08	4.9	High	Yes	Yes	C
220584167	13.8	2.1	15 ± 2	0.04	4.8	High	Yes	Yes	B
220596913	5.7	6.2	8 ± 1	0.09	3.7	High	Yes	Yes	B
910193711	6.0	1.8	30 ± 4	0.22	0.8	Low	No	Yes	C
910279515	3.5	1.5	23 ± 8	0.11	3.9	High	Yes	Yes	C

Notes. Column (2) gives the average S/N of the S/N map. The size of the galaxy with respect to the seeing is given in Col. (3). The criteria used to determine if galaxies are in rotation or not (8) are the mismatch between morphological and kinematic position angles (4) and the residuals normalized by the velocity shear (5). The ratio of the rotation velocity over the local velocity dispersion is given in Col. (6). Galaxies are classified according to the observed velocity shear in Col. (7): galaxies with $V_{\text{shear}} < 100 \text{ km s}^{-1}$ or $V_{\text{shear}} > 100 \text{ km s}^{-1}$ are respectively classified as low and high shear. The close environment classification and its associated quality flag are given in (9) and (10). The suffix “s” refers to secondary objects which have been detected both in $H\alpha$ and in the I -band image. ⁽¹⁾ This galaxy is classified as non-rotating even if it fulfils the criteria because the kinematics seem to indicate that this system is composed of two close companions (cf. Appendix A).

Usually they are quite young (~ 1 Gyr) but they can also be as old as ~ 10 Gyr, which would be compatible with our observations.

The intermediate mass galaxies ($\log M_{\text{star}} < 10.5$) are almost equally divided between rotating and non-rotating systems but the smallest ones are principally non-rotating systems

(see Fig. 4). Since both rotating and non-rotating systems are observed with small masses and radii, we can tentatively conclude that non-rotating systems are not caused by observational artifacts. However, we cannot exclude the possibility that for some of the latter systems, when $R_{\text{last}}/\text{Seeing} \sim 0.5$, the absence of

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Table 5. Physical properties of MASSIV subclasses.

Class	N	M_{star} [$10^{10} M_{\odot}$]			SFR [$M_{\odot} \text{ yr}^{-1}$]			R_e [kpc]			σ [km s^{-1}]		
		Med	Mean	Dev	Med	Mean	Dev	Med	Mean	Dev	Med	Mean	Dev
Rotating	19	3.9	4.0	3.7	46	60	44	3.1	3.8	2.4	62	58	19
Non-Rotating	16	1.2	1.6	1.2	30	39	29	2.7	2.7	1.2	60	61	17
Detected but unclassified	10	1.5	2.3	2.1	24	34	30	2.3	2.3	1.0	43	39	24
Isolated	28	1.4	3.0	3.4	37	47	41	2.7	3.1	2.2	62	56	22
Interacting	13	1.7	2.1	1.4	30	43	32	4.0	3.9	1.4	53	52	21
Rotating + Isolated	14	3.9	4.1	4.3	58	63	49	2.9	3.8	2.7	62	57	19
Rotating + Interacting	4	4.6	3.6	1.4	64	55	33	5.3	4.2	1.7	73	61	25
Non-Rotating + Isolated	8	1.4	2.0	1.7	42	37	22	1.9	2.0	1.1	66	64	19
Non-Rotating + Interacting	7	1.2	1.3	0.5	26	35	33	3.4	3.5	0.7	53	56	15

Notes. N is the number of galaxies in each subclass; med: median value; mean: average; dev: standard deviation. Only the main objects are considered.

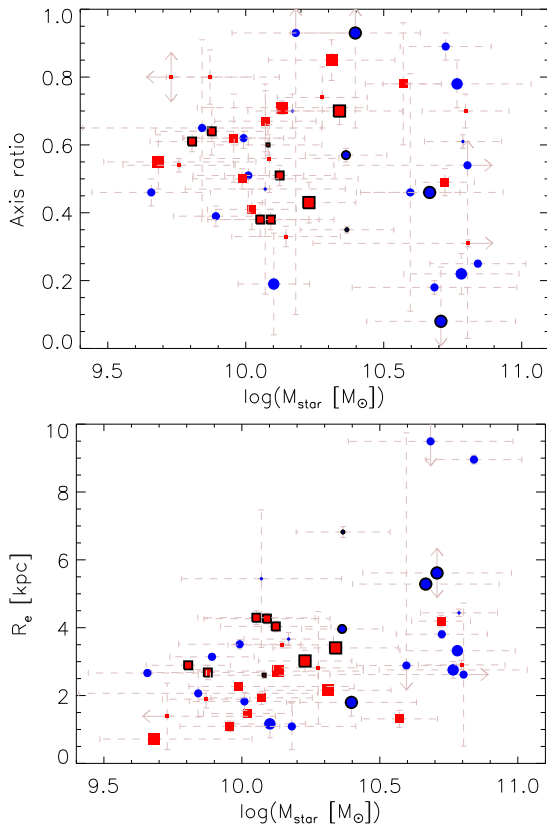


Fig. 4. *Top:* axis ratio as a function of the stellar mass. *Bottom:* half-light radius as a function of the stellar mass. Same symbols as Fig. 3. Arrows indicate that uncertainties are exceeding the displayed range.

observed rotation is due to an episode of star formation in single non-resolved regions. The fact that we observe large non-rotating galaxies incompatible with face-on systems is also very intriguing. In addition, those with $R_e > 2.8$ kpc all interact. This property is probably related to the nature itself of these systems.

5.2. Dynamical support of disks

To quantify the dynamical support, it is common to study the ratio of the rotation velocity over the local velocity dispersion. In

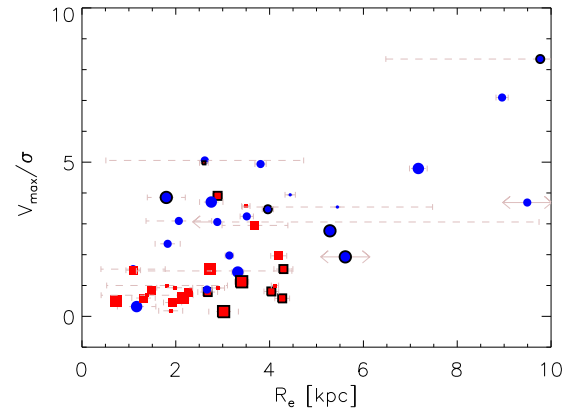


Fig. 5. V_{max}/σ as a function of the half light radius. Same symbols as Fig. 4.

Fig. 5, we show this ratio as a function of the half-light radius. Only two out of the 20 secure rotators in MASSIV have a ratio lower than unity and are consequently dispersion-dominated. Only five out of these 20 secure rotators have $V_{\text{max}}/\sigma < 2$, which indicates that the majority of rotating disks are clearly rotation-supported at $z < 1.5$. Moreover, there seems to exist a trend that the largest galaxies have the highest V_{max}/σ ratio. As seen in Table 5, the median velocity dispersion is $\sigma = 62 \text{ km s}^{-1}$ for rotators. In addition, the median rotation velocity $V_{\text{max}} = 141 \text{ km s}^{-1}$ and the median V_{max}/σ ratio is 3.2. By restricting the analysis on rotating systems with $R_e > 3.2$ kpc (median for all rotators), we end up with ten objects with a median velocity dispersion $\sigma = 49 \text{ km s}^{-1}$, a median rotation velocity $V = 201 \text{ km s}^{-1}$ and a median $V_{\text{max}}/\sigma = 3.7$, which clearly indicates that the largest disks are more stable. This seems to be supported in Fig. 6 where we clearly see that the velocity dispersion is low for the largest rotators ($R_e > 6$ kpc).

At higher redshift, rotators in the LSD/AMAZE sample ($z \sim 3.3$) have a typical gaseous velocity dispersion of $\sigma = 90 \text{ km s}^{-1}$ and $V_{\text{max}}/\sigma \sim 1.6$ (Gnerucci et al. 2011), whereas in the SINS sample ($z \sim 2.2$), Cresci et al. (2009) found $\sigma = 52 \text{ km s}^{-1}$ and $V_{\text{max}}/\sigma \sim 4.5$ as median values for their unambiguous rotators. The comparison with SINS is not straightforward since they have made these measurements only for a fraction of these systems (12 out of 18). For their sample of rotating systems, the median rotation velocity is $V_{\text{max}} \sim 240 \text{ km s}^{-1}$ and the median half-light radius (using the conversion between half-light

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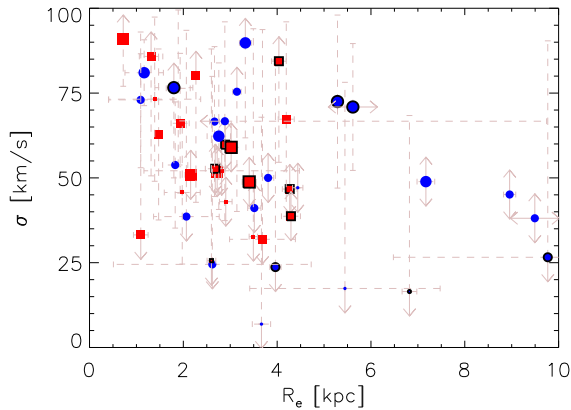


Fig. 6. Velocity dispersion as a function of the effective radius. Same symbols as Fig. 4.

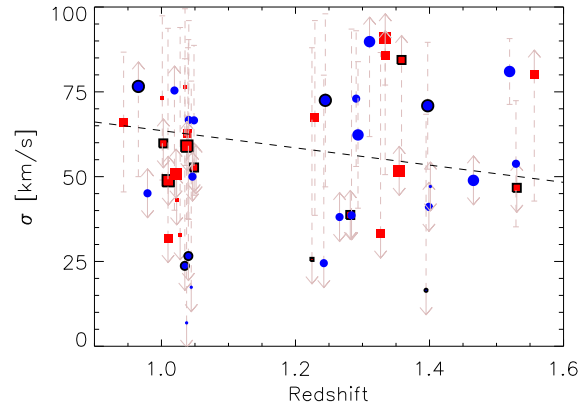


Fig. 7. Velocity dispersion as a function of the redshift. Same symbols as Fig. 4. The dotted line represents the velocity dispersion corresponding to the spectral PSF of SINFONI.

radius and disk scale length $R_e = 1.68R_d$ valid for an exponential disk distribution) is $R_e = 3.4$ kpc. These median values are higher than those we obtain for our sample of 20 rotating galaxies ($V_{\max} \sim 140$ km s $^{-1}$ and $R_e = 3.2$ kpc), which indicates that while their sample of rotation-dominated galaxies probes the massive disks population at $z \sim 2.2$, our sample of rotating systems spans a wider range in terms of dynamical mass. At lower redshift, from the IMAGES sample ($z \sim 0.6$), the typical gaseous velocity dispersion is $\sigma \sim 45$ km s $^{-1}$ (Puech, private communication) and $V_{\max}/\sigma \sim 4$ considering both rotating disks and perturbed rotators, with a trend for higher V_{\max}/σ for rotating disks (Puech et al. 2007). The GHASP sample is the largest 2D kinematics sample of local spiral galaxies with star-formation rates typical of the local Universe. We consider here the subsample of 136 galaxies that we projected at $z \sim 1.33$ with similar spatial resolution conditions as the MASSIV sample (see Sect. 3.2.3). These galaxies have a median optical radius of 8.5 kpc. Since $D_{25} \sim 1.9R_e$ for an exponential disk distribution, this gives a median half-light radius of 4.5 kpc. The median rotational velocity is 164 km s $^{-1}$, the median velocity dispersion is 24 km s $^{-1}$ and the median V_{\max}/σ is ~ 7 . Despite the small differences observed from $z \sim 2.2$ to $z \sim 0.6$, observations of various samples suggest a trend of decreasing velocity dispersion with decreasing redshift. Associated to local and $z \sim 3$ observations, this supports the idea that the gaseous phase of galaxy disks becomes less turbulent with cosmic time.

We also observe that about half of the sample of rotators have a gaseous velocity dispersion unambiguously higher than 60 km s $^{-1}$. These may be interpreted as clumpy disks. Indeed, it has been claimed that a high-velocity dispersion is expected for clumpy disks created from a smooth cold gas accretion (e.g. Bournaud & Elmegreen 2009). This mode of accretion is believed to be more efficient at $z > 2$ (Kereš et al. 2009) and it is therefore not surprising to also observe a significant percentage of galaxies in MASSIV that resemble stable rotating disks with velocity dispersions compatible with ~ 20 – 50 km s $^{-1}$ (cf. Fig. 7).

We also point out that the gaseous velocity dispersion is not significantly different when considering non-rotating systems, which raises the question of the origin of this gaseous velocity dispersion. If these non-rotating systems are indeed transient mergers of star-forming disks in an unstable phase, then a high-velocity dispersion is expected (Bournaud et al. 2011), but then it is more difficult to explain cases of large non-rotating galaxies with relatively low gaseous velocity dispersion.

As pointed out by e.g. Lehnert et al. (2009), Green et al. (2010), and Le Tiran et al. (2011), star formation might be responsible for the gaseous turbulence. In particular, star formation intensity might be correlated with gaseous velocity dispersion. Within this hypothesis, the lowering of the velocity dispersion from high to low redshift might be explained by the fact that the samples discussed in this paper have on average a decreasing star formation rate from $z \sim 3$ to $z = 0$ (see Contini et al. 2012). This behavior is due to cosmological surface brightness dimming but also to galaxy evolution itself: on average, galaxies have a larger size and form less stars at $z = 0$ than at $z \sim 2$. The relations between star formation and velocity dispersion will be addressed with the full MASSIV sample in a forthcoming paper.

5.3. Importance of merging at $z \sim 1.2$

The proportion of interacting galaxies in our sample is at least 29% (13 interacting systems with flag “A” or “B” out of the 45 detected systems with $0.9 < z < 1.6$) and is 32% if we only consider the 41 systems with a flag “A” or “B” in the redshift range $0.9 < z < 1.6$. Because of the way we determined which systems are in interaction, this proportion is a lower limit. Indeed, on the one hand, it could be that a fraction of non-rotating galaxies are ongoing mergers or merger remnants. On the other hand, close mergers could mimic rotators (elongation and velocity gradient along a common axis), as is the case for VVDS020294045 (see Appendix A). In addition, owing to the non-circular final field-of-view of our SINFONI data we only cover a fraction of the area at a given radius around our targets. There is also a substantial percentage of objects for which close galaxies are observed in *I*-band but not in SINFONI datacubes (see Appendix A). It is possible that a fraction of these objects are passive galaxies at the same redshift as the MASSIV targets. It is interesting to notice that on average, galaxies in interaction are larger than isolated galaxies and have a lower gaseous velocity dispersion (see Table 5).

The proportion of interacting systems in MASSIV is comparable to the percentage of mergers in the SINS sample. However, SINS mergers (around one third of the sample) are mainly identified using kinemetry, a technique based on the degree of perturbation observed in the kinematic maps, whereas the 29% of interacting galaxies in MASSIV are mainly based on the detection of several components. On the one hand, only three pairs out

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of 52 detected systems were identified in SINS (and three pairs out of 29 systems in the LSD/AMAZE sample), which is much lower than the pair fraction observed in MASSIV. On the other hand, SINS mergers found from kinemetry may mainly correspond to our non-rotating galaxies (see Sect. 4.4). That means that using this latter method, we would probably find that some MASSIV isolated galaxies are likely mergers. This is additionally supported by the fact that the percentage of non-rotating systems is larger in interacting galaxies: among the 11 systems with interactions in the sample (with both $S/N > 5$ and isolation flag “A” or “B”), seven have at least one non-rotating component (64%), whereas among the 22 isolated systems (with both $S/N > 5$ and isolation flag “A” or “B”), there are only eight non-rotating galaxies (36%). This could indicate that some isolated non-rotating galaxies are in fact ongoing mergers or merger remnants. This seems to be also supported by the number of interacting galaxies, which is larger at $z \sim 1$ (e.g. Fig. 7). The conclusion is therefore that the percentage of mergers at $z \sim 1.2$ is higher than at $z \sim 2.2$ as observed in SINS.

At lower redshift, 44% of galaxies in the IMAGES sample have complex kinematics that could be explained by ongoing mergers of star-forming galaxies (Yang et al. 2008). This is more than the proportion of interacting galaxies in MASSIV. However, the definition of galaxies with complex kinematics is probably closer to our definition of non-rotating galaxies (as is the case for the comparison with SINS) and the proportions of these two classes in each sample are fully comparable. It is probable that for IMAGES, a larger field of view would have led to the detection of close companions. Consequently, a direct comparison of merging with MASSIV is not straightforward.

The number of mergers deduced from the pair number is higher in MASSIV than in SINS. Our results are still consistent with a peak in the merger activity at $1 < z < 2$ if we consider systems with complex kinematics. However, it is not yet clear if these kinematic signatures of mergers are related to minor or major mergers, whereas from the MASSIV sample we have a way to infer the mass ratio between pair members. A dedicated analysis on the merger rate from the observed pairs will be performed for the entire MASSIV sample (López-Sanjuán et al., in prep.) and will probably enable us to infer the rate of observed ongoing mergers.

5.4. Nature of non-rotating galaxies

Clearly, non-rotating systems are mainly galaxies classified as low-velocity shear galaxies plus some interacting galaxies. Among the 16 (and eight additional with $S/N < 5$) low-velocity shear galaxies only three (and two additional that have very poor S/N) could be considered as rotating. Two of these galaxies are compatible with nearly face-on systems. The exact nature of these non-rotating objects is still unclear. Such a population of galaxies has already been observed at higher redshift both in the SINS sample at $z \sim 2.2$ and in the LSD/AMAZE sample at $z \sim 3.3$. These objects are smaller on average than rotators and are often associated to interacting systems (see Table 5).

An unexpected trend is observed for these objects in the MASSIV sample (Fig. 6): there exists an anti-correlation between the mean velocity dispersion of the gas and the effective radius of the stellar component. This correlation is in contrast to what would be expected for elliptical galaxies located in the fundamental plane (Dressler et al. 1987; Djorgovski & Davis 1987). However, the fundamental plane applies for the central stellar velocity dispersion, which is a quantity that cannot currently be measured in high-redshift galaxies from spectroscopic data.

Alternatively, this correlation could reflect a beam-smearing effect since an unresolved velocity gradient could reproduce this trend. However, the fact that we also observe velocity gradients in other galaxies with similar sizes contradicts this interpretation but it is closely linked to the spatial distribution of the ionized gas. For face-on disks (possible only for low-velocity shear galaxies), it would be expected to recover the same trend as for rotators (i.e. lower velocity dispersion for large disks). This is very unlikely, however, because velocity gradients may be observed for the largest galaxies and also because we do not expect to observe more than four galaxies with an inclination lower than 20° (see Fig. 2). If the gas were more concentrated, though, it might collapse more violently to form stars at a higher density if we consider that star formation can drive the gaseous velocity dispersion as suggested by e.g. Green et al. (2010).

As was been proposed in the previous sections, these non-rotating objects can be interpreted as

- Merger remnants in a transient unstable state for the gaseous phase (e.g. collapse of gas due to interaction, anti-spin mergers, etc.).
- Rotators with a special gaseous distribution: the most plausible being a unique highly concentrated star-forming region when galaxies are small.
- Galaxies with an unstable gaseous phase. This could be supported by the relatively low mass (thus a low potential well) of these systems.
- Nearly face-on disks (for low-velocity shear galaxies only): this would explain the similar velocity dispersion as for the rotators but the number of these galaxies is statistically low.
- Spheroids: however, (i) these galaxies may be the most massive ones; (ii) there is no reason to think that the gas is not in a disk as observed at lower redshift (e.g. Peletier et al. 2007 in early-type spirals; e.g. Bournaud et al. 2007 in numerical simulations of gas-rich mergers) and (iii) if the gas is randomly distributed, a higher velocity dispersion than for rotators would be expected.

6. Summary and conclusions

We presented the data and their processing for a subsample of 50 galaxies of the MASSIV sample (described in Contini et al. 2012). We focused on the analysis of the dynamical state of galaxies using kinematic maps derived from SINFONI data-cubes. We classified this sample based on the modeling of these maps and based on a comparison between morphology and kinematics. This classification describes on the one hand the rotating or non-rotating nature of the objects and on the other hand the close environment of galaxies.

We separated rotating galaxies and non-rotating galaxies based on two arguments: (i) the agreement between morphological and kinematic major axis position angle and (ii) the accuracy of the description of a rotating disk model. We showed that about half of the sample displays rotation, whereas one third does not show a dominant ordered rotation (the remaining fraction is not classified). However, the ionized gas turbulence in these two classes is similar ($\sigma \sim 60 \text{ km s}^{-1}$), which marks a transition between higher redshift galaxies showing on average a higher velocity dispersion ($\sigma \sim 60\text{--}90 \text{ km s}^{-1}$), as in SINS at $z \sim 2.2$ or LSD/AMAZE at $z \sim 3.3$, and lower galaxies in IMAGES at $z \sim 0.6$ and GHASP at $z = 0$, which are characterized by a lower velocity dispersion ($\sigma \sim 20\text{--}40 \text{ km s}^{-1}$). All these samples are 3D spectroscopic samples observing ionized gas emission lines. This transition of the gaseous velocity dispersion seen

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from these high- to low-redshift samples is partly responsible for the increase of the rotational support (deduced from the V_{\max}/σ ratio) when galaxies evolve. It could be related to a common process that would induce a decrease of the star formation rate. We observe that about half of our galaxies have a velocity dispersion compatible with that in local star-forming galaxies, whereas the other half is more compatible with galaxies at higher redshift. At high redshift, this high-velocity dispersion seen in isolated disks is thought to be caused by a mass assembly driven by cold flows. In this framework, our results suggest that at $z \sim 1.2$ cold gas accretion is less efficient than at higher redshift but more efficient than at lower redshift, consistent with cosmological simulations (Kereš et al. 2005). When we compare our sample to these other high-redshift samples, it seems that the fraction of disks increases in star-forming galaxies while the Universe evolves.

By studying strong kinematic signatures of merging and detecting pairs in our data-cubes and broad band images, we have shown that the fraction of interacting galaxies is up to at least one third of the sample. The fraction of non-rotating objects in these systems is higher than in isolated ones. This suggests that a significant fraction of isolated non-rotating objects could be mergers in a transient state in which the gas is not dynamically stable. However, the nature of non-rotators is still unclear. Compared to higher and lower redshift 3D spectroscopic surveys, our findings seem to indicate that at the peak of star formation activity, the fraction of star-forming galaxies in interaction is also at its maximum, corresponding to a peak in the merging activity (Ryan et al. 2008; Conselice et al. 2008; López-Sanjuan et al. 2009).

All these results favor a scenario in which the mass assembly of star-forming galaxies is progressively shifting from a predominance of smooth cold gas accretion to a predominance of merging as cosmic time evolves, with a transition epoch around a redshift $z \sim 1-1.5$.

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Appendix A: Individual comments

VVDS020106882 This galaxy has a regular velocity field. The velocity dispersion is higher on the southeast side but remains low. Some external pixels are detected on the northwest side. They seem to be related to regions associated to the galaxy due to concordant velocities. Two objects (including a very faint one) are detected in *I*-band at less than 6'' but are not detected in the SINFONI datacube.

VVDS020116027 This galaxy has a small companion at 4'' to the north. The companion is detected in H α but with a low *S/N* and over a region of the size of the seeing. This companion is less exposed than the main galaxy. The main galaxy itself seems to have two components in the H α flux distribution and has a very low velocity shear.

VVDS020147106 Despite a low-velocity gradient and some perturbations in the low *S/N* regions, the kinematic position angle of this galaxy agrees well with the *I*-band morphology. This galaxy may therefore be a nearly face-on disk (at least the ionized gas) but with a high velocity dispersion. There are several objects detected in *I*-band but not in the SINFONI datacube at less than 6''.

VVDS020149061 The *I*-band image suggests small and extremely faint companions that are not detected in the SINFONI datacube, however. The seeing is half the size of the H α emission. The velocity field is irregular on the edges where the *S/N* is lower and the velocity dispersion is higher than 70 km s⁻¹ everywhere in the galaxy (except at the edges).

VVDS020164388 Except for the edge, the velocity field and the velocity dispersion map are regular. Velocity dispersion is not higher than 60 km s⁻¹. In the *I*-band image, the galaxy is quite round. Several faint close objects are also observed in this image but are not detected in H α .

VVDS020167131 Some [OIII] is detected in the center but there are only a few pixels and the *S/N* is lower than 4. In addition, some emission line is detected at the same redshift in the center of the galaxy on the east seen in the *I*-band. This indicates that these two galaxy are in interaction.

VVDS020182331 The velocity gradient is clear even if we exclude low *S/N* regions (bluest and reddest velocities). However, the velocity field and velocity dispersion map look perturbed, probably because of the low *S/N* of the observation. A very close companion is detected in the *I*-band image that is undetected in H α .

VVDS020193070 Despite a generally low *S/N*, the velocity field is quite regular except in the outskirts. The velocity dispersion is low but slightly asymmetric. The morphology is elongated both in *I*-band and in H α (but with a difference of $\sim 10^\circ$ between morphological and kinematic major axis position angles). Three very faint objects are detected at less than 5'' from the galaxy but are not detected in H α .

VVDS020208482 The *S/N* is fairly low but there is a clear velocity shear. The *I*-band image shows a close galaxy and other objects, farther away, that are not detected in H α .

VVDS020214655 This galaxy displays a fairly smooth low velocity gradient, but not aligned with the morphology, which is quite round. There are some perturbations in the low *S/N* regions of the velocity field. There are also small perturbations in the velocity dispersion map, which is fairly flat and has low

values. There is a close galaxy in the *I*-band (at $\sim 2''$) that is not detected in H α .

VVDS020239133 The morphology from *I*-band is clearly elongated and is more extended than that in H α . The velocity field is asymmetric but shows a clear gradient. The velocity dispersion map is also perturbed and has a high-velocity dispersion in the south that could be explained by a sky line residual. Two objects are observed in *I*-band but not detected in H α within the SINFONI field-of-view.

VVDS020240675 The velocity field is perturbed, with a low-velocity gradient. The velocity dispersion is fairly low (~ 40 km s⁻¹). In addition, the size of the H α emission is not large compared to the seeing. There is some emission detected at 2'' in the east of the main galaxy, but nothing is detected in the *I*-band at this location. There are also several objects in *I*-band within 6'' that are not detected in H α .

VVDS020255799 The signal in H α for this galaxy is not extended and with a mean *S/N* lower than 5. There is no clear regular velocity gradient. One object is detected at 5'' in *I*-band but not in H α .

VVDS020261328 The velocity field is not completely regular in the blue and red sides. There is also a peak in the velocity dispersion where the *S/N* is lower. Several objects are detected in *I*-band but not in H α .

VVDS020278667 Some H α is detected, but the extent is lower than the seeing, and the *S/N* is lower than 5. Therefore its classification is very doubtful. There is a close galaxy ($\sim 1.5''$) in the *I*-band but it is not detected in H α .

VVDS020283083 This galaxy looks asymmetric in the *I*-band and this coincides with the H α distribution and with the main perturbation of the velocity field. This could be related to a tidal tail. That is why this galaxy is classified as non-isolated.

VVDS020283830 In the H α flux map there is a faint detection smaller than the seeing that coincides with a detection in the *I*-band (north-east). For the main object, the distribution is asymmetric. This is also true in the H α distribution, which is clumpy. These clumps may be either star-forming regions in one single rotating galaxy or two small objects in the process of merging.

VVDS020294045 The total velocity shear is larger than 50 km s⁻¹. However, both the *I*-band and the H α images are consistent with a system involving two components: one main component with a low-velocity gradient (south) and one small component with a size comparable to the seeing (north). This interpretation is supported by the high-velocity dispersion between the two components due to beam smearing between the two components at different radial velocities. There is a sky line residual at the velocity of the small component, but the flux detected is noticeable. It is therefore not possible to classify the small component. This system is classified as interacting with a flag B since it could also be a single object but with strong perturbations.

VVDS020363717 The *I*-band shows a close ($\sim 1''$), very diffuse and faint object that is not detected in the H α (southwest). The velocity shear is low. The velocity field is perturbed in the low *S/N* regions. The velocity dispersion map is also perturbed and the velocity dispersion is high (~ 90 km s⁻¹) after beam smearing correction.

VVDS020370467 Both velocity field and velocity dispersion map are perturbed. The velocity gradient of the model is low

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and the velocity dispersion is quite high (higher than 70 km s^{-1}) in particular at the north edge (higher than 150 km s^{-1}). This could be a sign for a strong interaction due to merging, but there is also a sky line residual that could induce the line width. In addition some objects are observed in the *I*-band image at less than $6''$ but are not detected in $H\alpha$.

VVDS020386743 The $H\alpha$ distribution is very elongated for this galaxy, but no velocity gradient is visible. Both velocity field and velocity dispersion are smooth. The *I*-band image shows a distorted galaxy compatible with the $H\alpha$ distribution. In the north, the distortion is compatible with a faint detection in $H\alpha$ that could be related to a minor close companion.

VVDS020461235 The velocity field of this galaxy is perturbed. There is a blob detected in $H\alpha$ in the southwest side. This blob is also suggested in the *I*-band morphology but is less clear. This blob has velocities compatible with the rotation of the main component but could as well be a minor companion in the process of merging. This system is therefore classified as interacting with a flag B.

VVDS020461893 The velocity field is regular but the velocity dispersion is less regular and high and could suggest a high-dispersion clump. An object is observed at $3''$ southward in the *I*-band image but is not detected in SINFONI data.

VVDS020465775 Both velocity field and velocity dispersion map are irregular. The gradient of the model is lower than 50 km s^{-1} . We claim that there is a companion in the northwest that induces broad and non-Gaussian (hence a lower *S/N*) lines interpreted as a sign of interaction that could also be responsible for the asymmetry in the *I*-band. Hence it is classified as interacting with a flag B since it could also be a single object but with strong perturbations.

VVDS140083410 The velocity field is perturbed and there is no clear overall velocity gradient. No galaxy is detected at less than $8''$ in the *I*-band image.

VVDS140096645 The velocity field is smooth but with some perturbations along the minor kinematic axis. The velocity dispersion map is perturbed and larger on the southern side probably due partially to a sky line residual. Some emission is detected in the $H\alpha$ map outside the main component (north). This detection is smaller than the seeing and is not clearly confirmed in the *I*-band morphology (CFHT12k images are less deep than CFHTLS images). It could be a minor companion but it is also compatible with the velocity of the main component and could be associated to it. Therefore it is classified as interacting with a flag B. This galaxy also has an AGN (see [Queyrel et al. 2012](#)).

VVDS140123568 Some $H\alpha$ is detected, but there are only a few pixels and the *S/N* is lower than 4. Therefore its classification is very doubtful. There is no detected object in the *I*-band image at less than $6''$ around the galaxy. Since the kinematics does not allow to investigate a possible ongoing merger, the isolation has a flag B.

VVDS140137235 Some $H\alpha$ is detected, but there are only a few pixels and the *S/N* is lower than 4. Therefore its classification is very doubtful. There is no detected object in the *I*-band image at less than $6''$ around the galaxy. Since the kinematics does not allow to investigate a possible ongoing merger, the isolation has a flag B.

VVDS140217425 The velocity field is regular and has the highest shear ($\sim 600 \text{ km s}^{-1}$). The velocity dispersion map shows two

regions with high-velocity dispersions ($>150 \text{ km s}^{-1}$). These regions are associated with a lower *S/N*. This is because there are double profiles in these regions that induce a poor fit. These double profiles are caused by the beam-smearing that mixes regions with different velocities. This can be interpreted as the presence of star-forming clumps at the edge of the galaxy and a large clump in the center. The decomposition of the profile using two Gaussian scales down the velocity dispersion to usual values, and shows that the two external clumps have no velocity gradient, whereas the central one does have one. This picture is compatible with a rotating disk in which the rotational velocity reaches a plateau before the external clumps. Using the double profile decomposition instead of the kinematics model to derive the maximum velocity, we find $V_{\text{max}} \sim 312 \pm \sim 20 \text{ km s}^{-1}$, which is much more realistic than the value found from the model. The velocities of the plateau are $\sim +313$ and ~ -312 , which shows that the rotation curve is fairly symmetric and therefore favors a rotating disk hypothesis. Correcting for the inclination, we obtain $V_{\text{max}} = 322 \text{ km s}^{-1}$. However, the *I*-band image is asymmetric (which is not incompatible with the previous statement), and it cannot be ruled out that the external clumps are instead objects about to merge with the main component.

VVDS140258511 The velocity field and velocity dispersion map look perturbed at the northeast edge but this may be due to a sky line residual. Except at this location, the velocity dispersion is low (less than 40 km s^{-1}).

VVDS140262766 The velocity field and the velocity dispersion map are not strongly perturbed and the velocity dispersion is about 50 km s^{-1} . Several objects are seen at less than $5''$ in the *I*-band image but are not detected in $H\alpha$.

VVDS140545062 This object has a clear and smooth velocity shear compatible with the *I*-band and $H\alpha$ morphology. However, the velocity field is not very well reproduced by the rotating disk model: one side is faster than the other. The velocity dispersion map shows a peak in the southwest side. This peak is caused by a double profile that may be interpreted as a sign of merging (late stage or ongoing merger). However, since the system as a whole is isolated (no galaxy at less than $7''$ in *I*-band), this galaxy was classified as isolated with a flag B.

VVDS220014252 Both velocity field and velocity dispersion map are perturbed. In addition, the morphology in *I*-band is asymmetric. This system has a high-velocity dispersion ($>80 \text{ km s}^{-1}$) and is peaked on some edges. This may be a sign of merging (ongoing or remnant). However, the system is isolated and there is no other evidence for some companion except one very faint object in *I*-band at $\sim 5''$, which is not in the SINFONI field-of-view. So this system is classified as isolated.

VVDS220015726 The velocity field of this galaxy is well reproduced by a rotating disk model. The velocity dispersion map is peaked in the center, which agrees with the expected effect of beam smearing caused by the inner velocity gradient. The velocity field and velocity dispersion map are slightly perturbed. One very faint component is detected in *I*-band at around $4''$ but is not detected in $H\alpha$.

VVDS220148046 This galaxy has been observed using AO and is therefore less deep than seeing-limited observations. Some emission line was detected but with a low *S/N* (~ 4). It appears that the redshift determined from the VVDS was wrong and that we observed [OIII] at $z = 2.2442$. The emission is compact and therefore it is not possible to conclude about the kinematics.

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VVDS220376206 A small component (smaller than the seeing, hence a flag B for the environment classification) is detected in $H\alpha$ and is confirmed in the I -band image (north). The residual velocity field shows signs of interaction. The main component has two blobs in $H\alpha$ and the I -band morphology is slightly asymmetric. The velocity dispersion map is also perturbed and peaked on the edges. This may be a sign of interaction or of the end of a merging event.

VVDS220386469 This observation was made with AO and is underexposed (owing to the small pixel size, cf. [Contini et al. 2012](#)). Even if the AO observation does not allow a fair comparison with other galaxies in the same conditions because of the smaller field of view, this system is classified as isolated with a flag B since the closest clear detection in the I -band image is farther than the $6''$ that would be reached using the seeing-limited observing strategy. The velocity field is irregular and the velocity dispersion is very small but this may be due to the use of AO, which provides more details and a lower S/N than seeing-limited observations. The extent of the velocity field is small so it is difficult to argue that there is no velocity gradient at all in this galaxy.

VVDS220397579 This system is the clearest interacting system between two large galaxies. The two objects are detected in $H\alpha$ and in the I -band. Both are elongated toward the same direction. The main object in $H\alpha$ and in I -band has a smooth velocity field with no gradient except on the northern side, near the companion. This region coincides with a higher velocity dispersion and this is probably related to the interaction with the companion. Since the galaxies are aligned and since the overall velocity gradient is monotonic, the whole system might be considered as a unique galaxy with two big clumps. However, such a system would be very much extended, a property which is physically unlikely at these redshifts. The companion shows a clear velocity shear and has been classified as well. It is also worth noticing that the minor companion in I -band is the most luminous in the K -band (from UKIDSS Deep Extragalactic Survey, [Lawrence et al. 2007](#)), therefore likely the most massive one.

VVDS220544103 This galaxy has a clear velocity gradient. The object is distorted in the I -band image as well as in $H\alpha$. This can be explained by the presence of two components (the morphology was decomposed using this hypothesis). In that case, the southern component may be well described by a rotating galaxy and the brightest blob in $H\alpha$ (north) maybe a companion in a merging stage or a relic of merging. However, it cannot be ruled out that this is a single object with disturbed morphology and kinematics, hence the flag B.

VVDS220544394 The $H\alpha$ map clearly shows the presence of a main component and a faintest one in the north. Their velocity fields seem decorrelated enough to claim that these are two separated objects on a pre-merging stage. The small component is also clearly detected in the I -band. The velocity field of the main component is perturbed but shows a velocity shear. Its velocity dispersion is about 50 km s^{-1} . The small component is slightly larger than the seeing and may also be considered to have a velocity shear.

VVDS220576226 The velocity field is perturbed in the north where the S/N is the lowest. The velocity dispersion map is regular. Two objects are observed within $6''$ in the I -band image but are not detected in $H\alpha$.

VVDS220578040 The I -band image shows an asymmetric galaxy with a more diffuse emission in the east. The $H\alpha$ map shows two blobs. These can be interpreted either as two distinct galaxies or as two clumps in one single galaxy. The velocity field is perturbed and is not fully reproduced by the model. The velocity dispersion map is slightly perturbed. Since no strong kinematics perturbation is observed, this galaxy is classified as isolated (with a flag C) and therefore its kinematics is interpreted as a sign of rotation.

VVDS220584167 The I -band morphology is distorted and the $H\alpha$ distribution is asymmetric. The kinematic position angle, however, agrees very well with the morphology. The velocity field shows some perturbations such as a slight asymmetry. This could be due to the position of the kinematic center that would differ from the morphological one. The velocity dispersion map is peaked in the center, however, but the beam smearing may not account for the whole amplitude of the peak. Two objects are detected in I -band but not in $H\alpha$ at less than $6''$.

VVDS220596913 This galaxy is a chain galaxy in the I -band. It has been observed both with and without AO (cf. [Contini et al. 2012](#)). Several blobs are detected in $H\alpha$. They can be interpreted as several clumps in a single edge-on galaxy since the velocity field is monotonic. They can also be interpreted as several small galaxies in a merging stage. The continuum was extracted from the SINFONI data around $H\alpha$ and seems to be peaked in the center, between the $H\alpha$ blobs, which favors the edge-on hypothesis. Since no strong kinematics perturbation is observed, and since a previous seeing-limited observation ([Epinat et al. 2009](#)) revealed that no extra emission is detected, this galaxy is classified as isolated with a flag B and therefore its kinematics is interpreted as sign of rotation.

VVDS910193711 This observation was made with AO. The extent is larger than $1''$. The velocity field is perturbed. The interpretation of the perturbations may not be comparable with seeing-limited observations because the S/N is lower and the scale of the perturbations smaller. The velocity dispersion is high (around 100 km s^{-1}). In the close neighborhood that would be observed using the seeing-limited strategy, several objects are detected in the I -band image.

VVDS910279515 This galaxy was observed using AO and is therefore less deep than seeing-limited observations. This galaxy is unclassified since there are only a few pixels with a S/N larger than 3. There are several objects in the I -band image at less than $5''$ that would have been observed using a seeing-limited strategy.

Appendix B: Kinematic maps

This appendix shows the kinematic maps for all detected galaxies. There are several sets of maps for galaxies with detected companions: one for the whole system (without kinematics modeling) and one for each modeled component.

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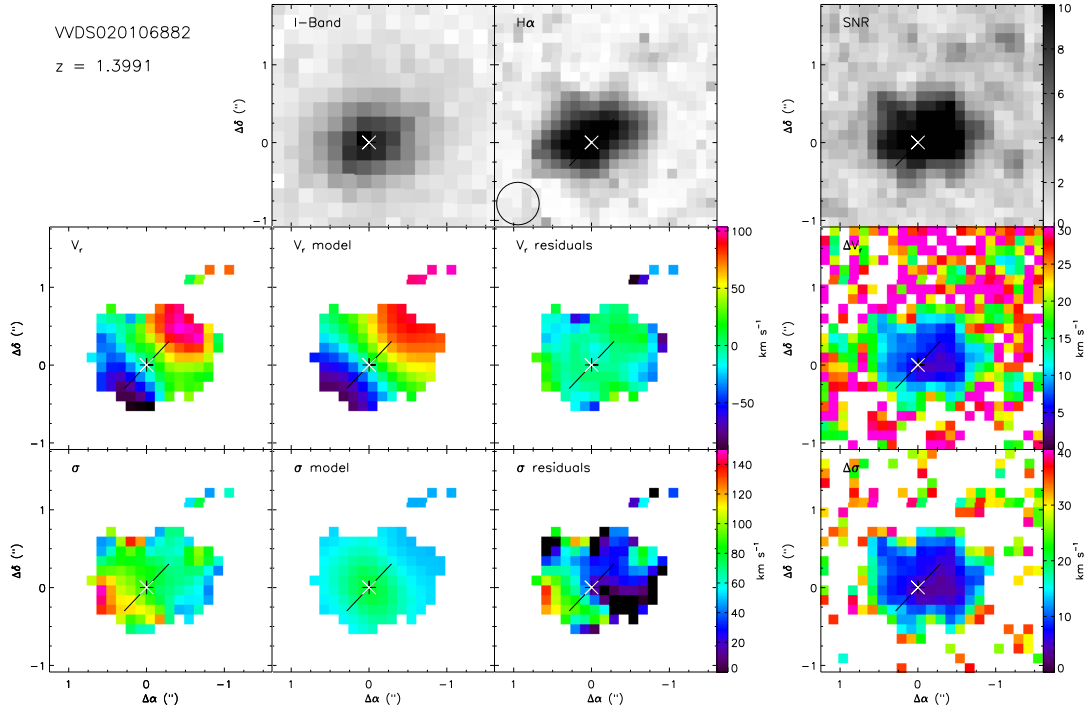


Fig. B.1. Maps for VVDS020106882. *From left to right: (top) the I-band CFHT image (arbitrary scale), the H α flux map (arbitrary scale) and the S/N map, (middle) the observed velocity field, the rotating-disk-modeled velocity field, the residual velocity field, the uncertainty map on the velocity field, (bottom) the observed uncorrected velocity dispersion, the velocity dispersion map deduced from the velocity field model (beam-smearing effect and spectral PSF), the beam-smearing-corrected velocity dispersion map and the uncertainty on the velocity dispersion. The redshift is indicated in the top-left. In each map, north is up and east is left. The center used for kinematics modeling is indicated as a double black and white cross, the position angle is indicated by the black line. This line ends at the effective radius. The seeing FWHM is indicated on the H α map as a circle.*

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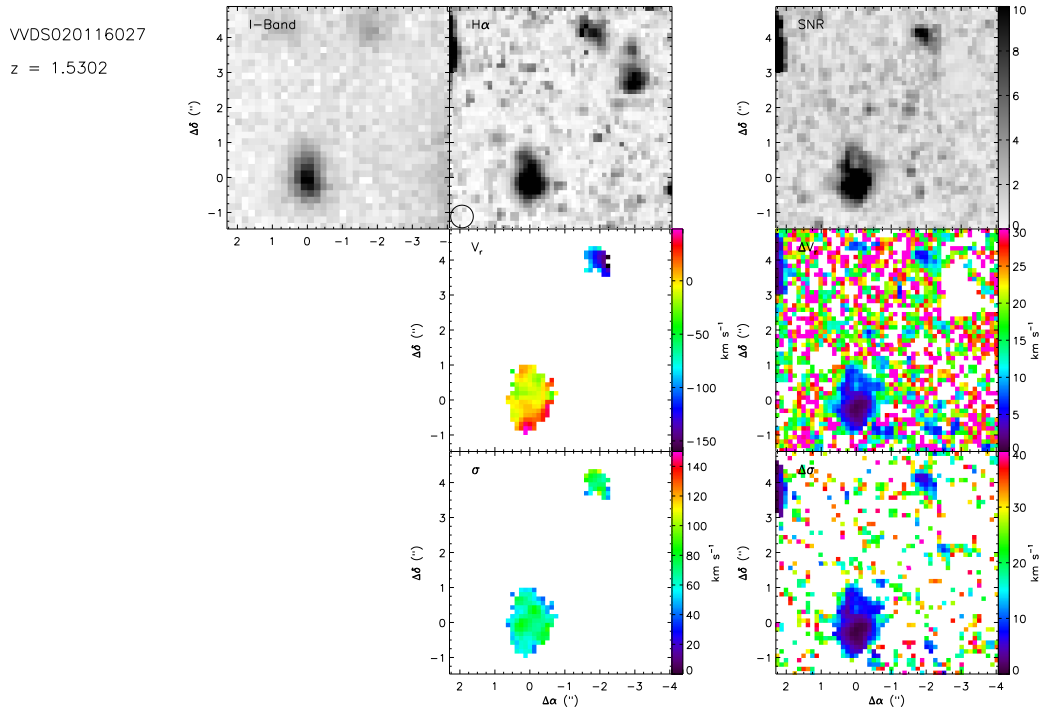


Fig. B.2. Maps for VVDS020116027 system. From *left to right*: (*top*) the *I*-band CFHT image (arbitrary scale), the *H* α flux map (arbitrary scale) and the *S/N* map, (*middle*) the observed velocity field, the uncertainty map on the velocity field, (*bottom*) the observed uncorrected velocity dispersion and the uncertainty on the velocity dispersion. The redshift is indicated in the *top-left*. In *each* map, north is up and east is left. The seeing FWHM is indicated on the *H* α map as a circle.

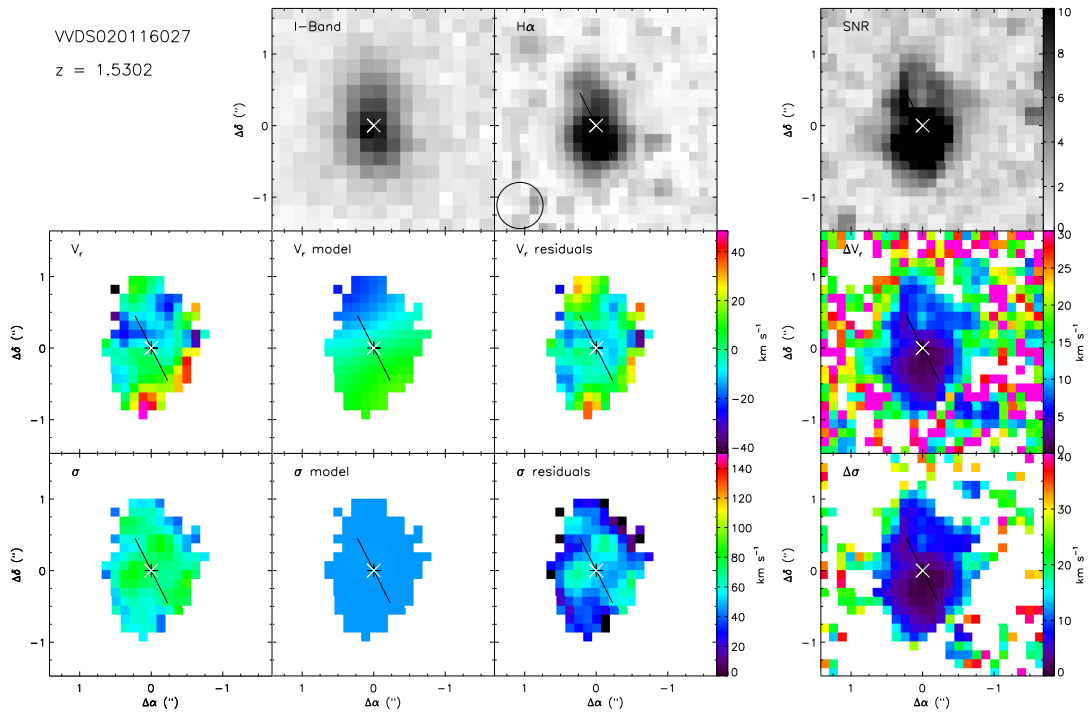


Fig. B.3. Maps for VVDS020116027. Same caption as Fig. B.1.

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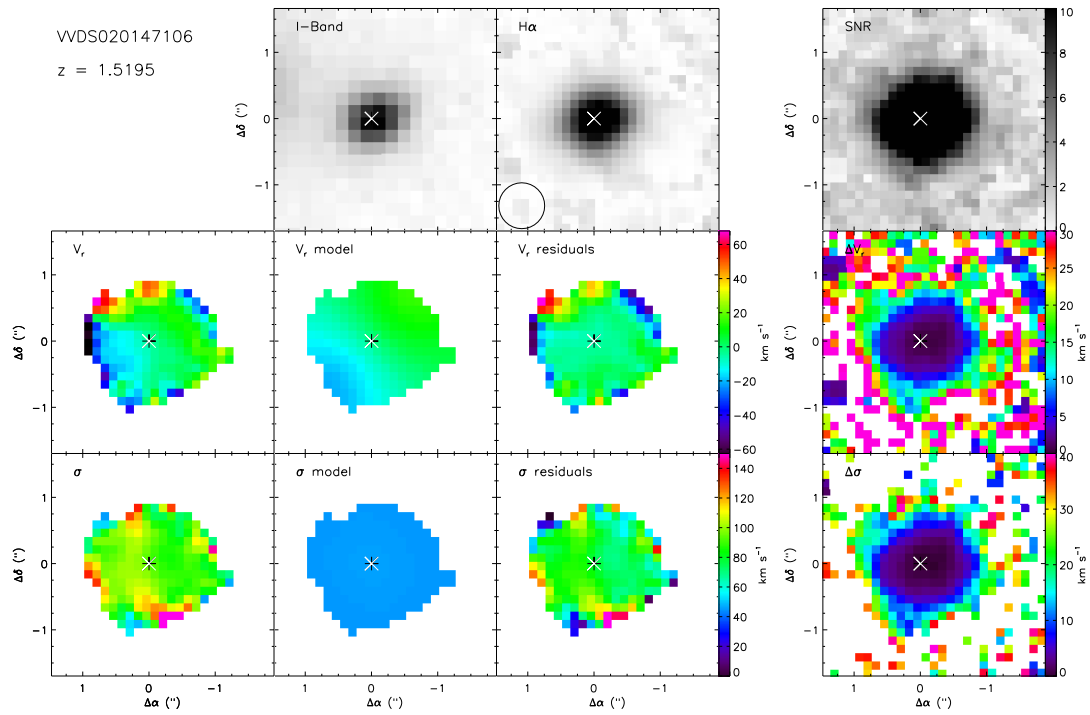


Fig. B.4. Maps for VVDS020147106. Same caption as Fig. B.1.

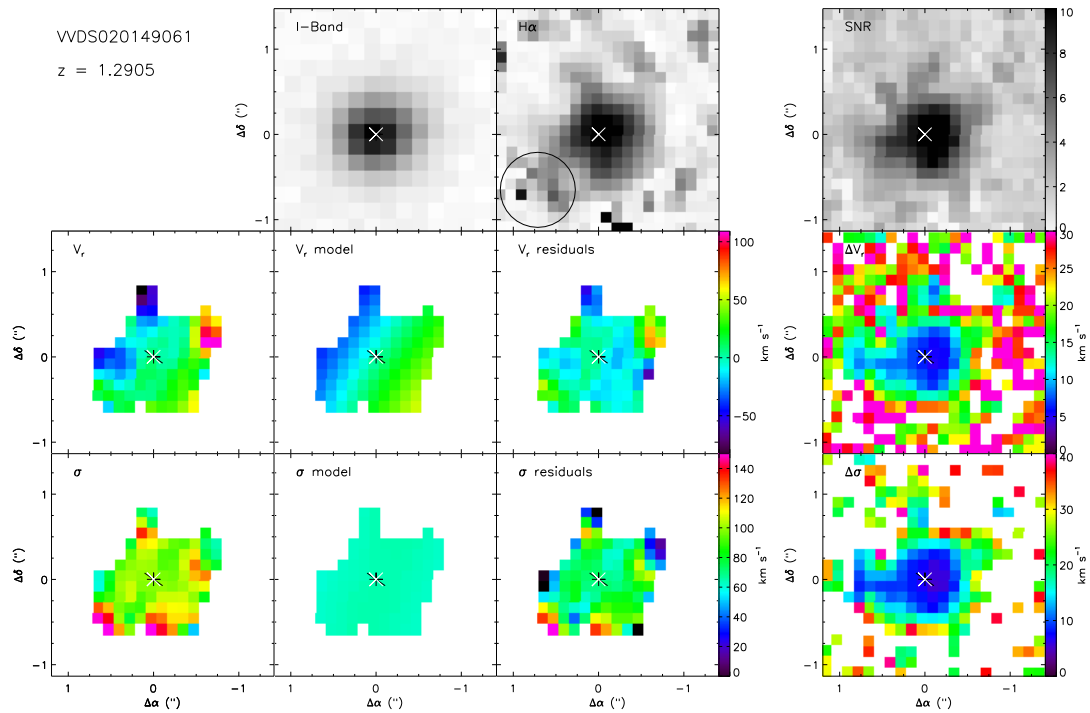


Fig. B.5. Maps for VVDS020149061. Same caption as Fig. B.1.

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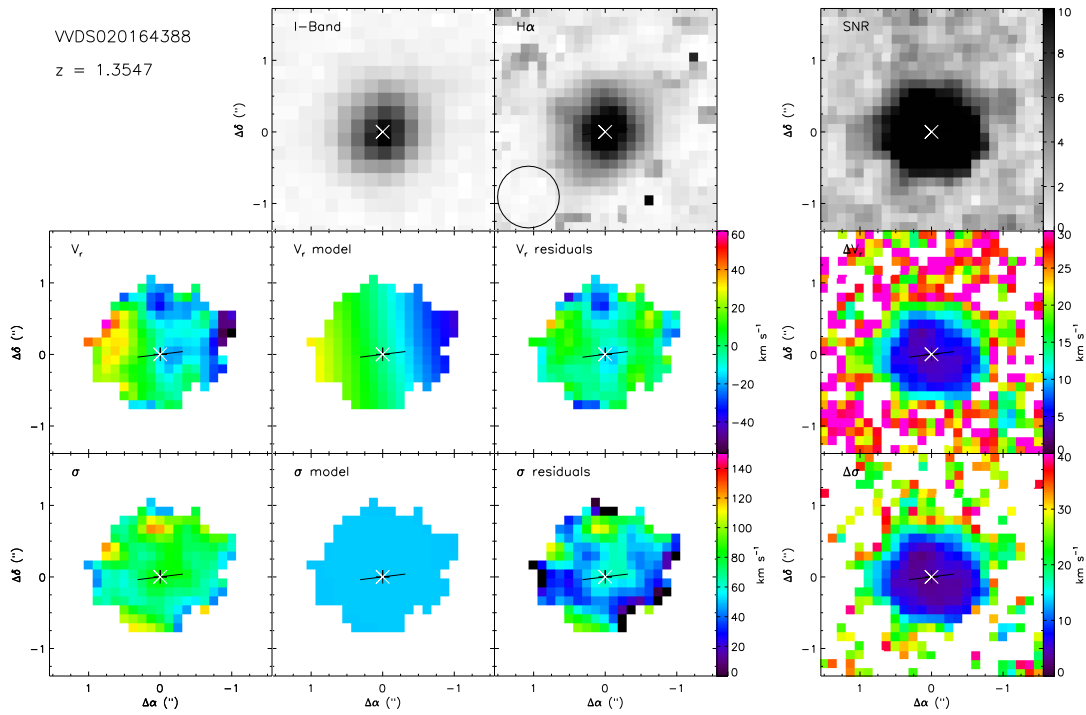


Fig. B.6. Maps for WVDS020164388. Same caption as Fig. B.1.

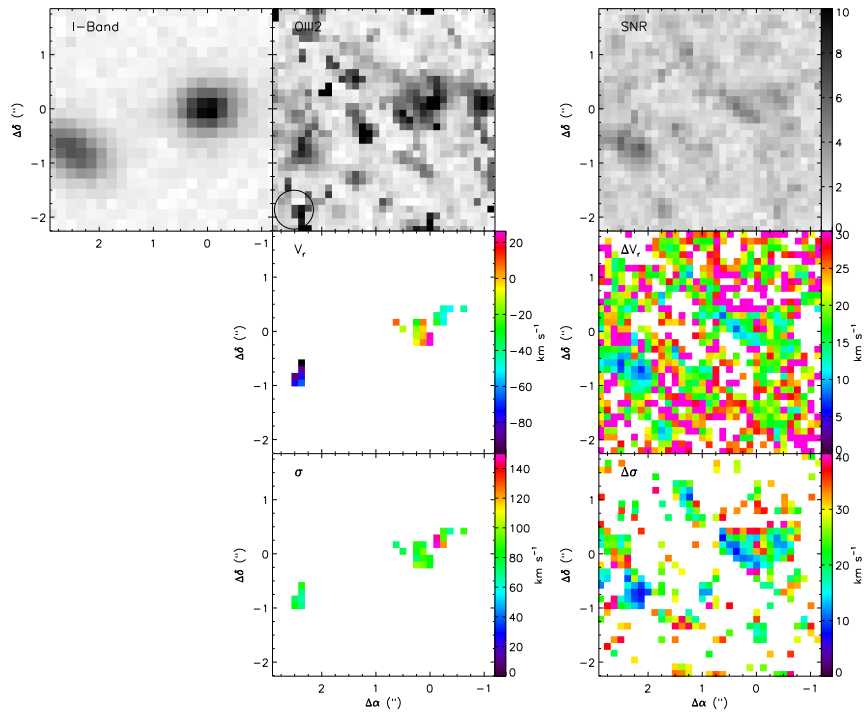


Fig. B.7. Maps for WVDS020167131 system. Same caption as Fig. B.2.

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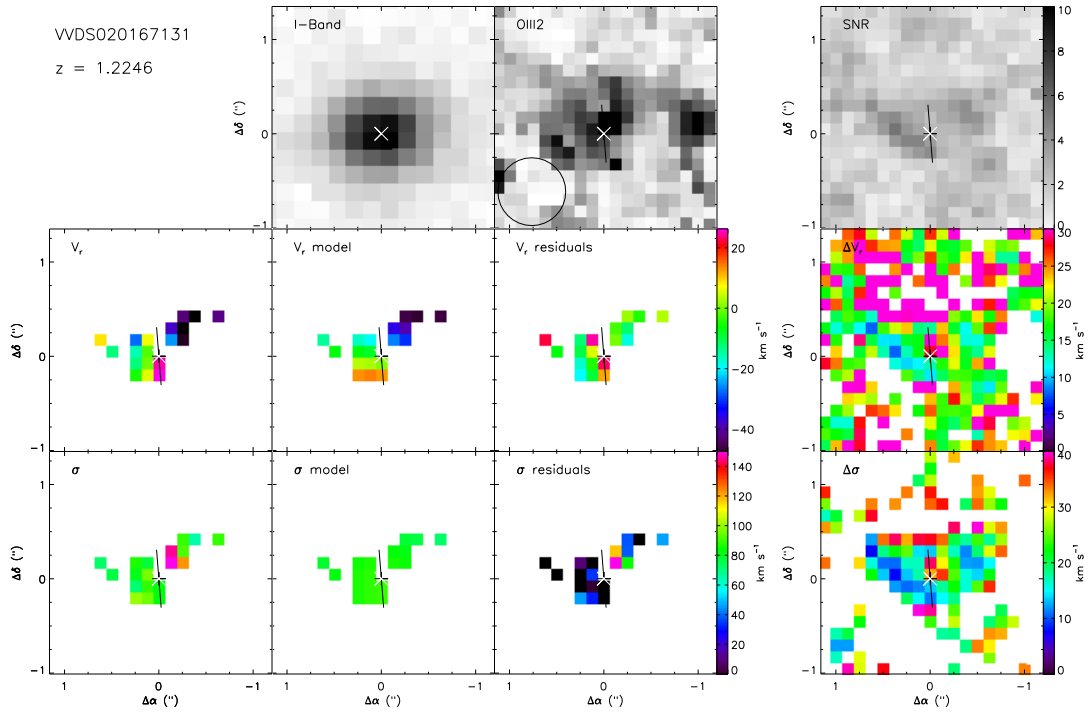


Fig. B.8. Maps for VVDS020167131. Same caption as Fig. B.1.

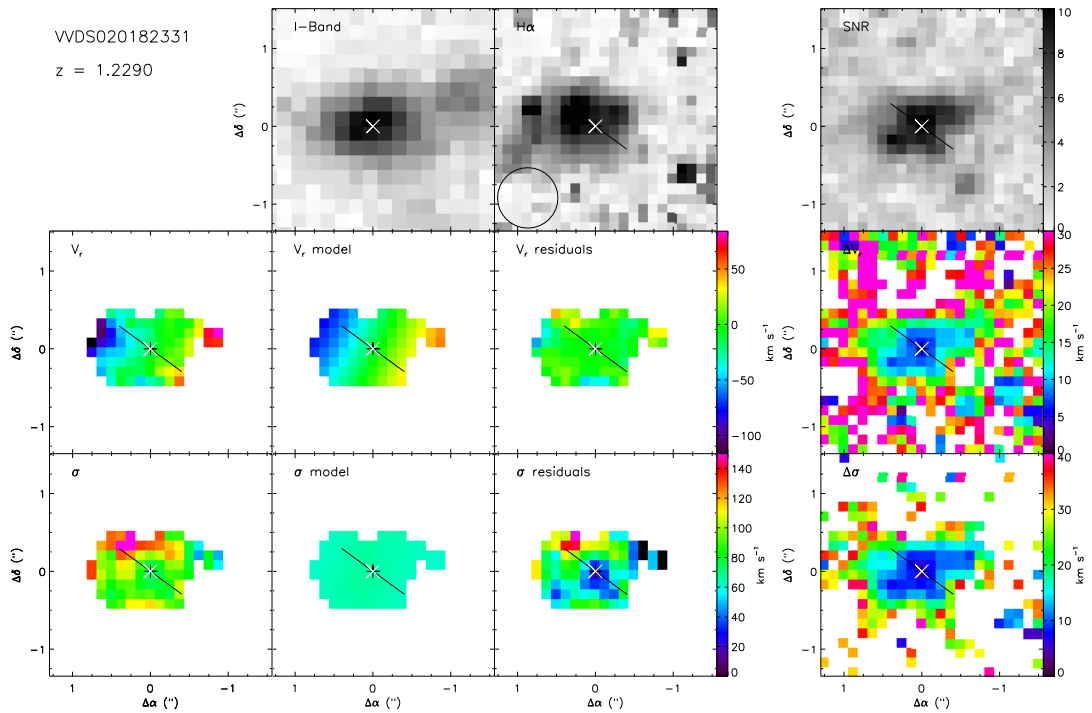


Fig. B.9. Maps for VVDS020182331. Same caption as Fig. B.1.

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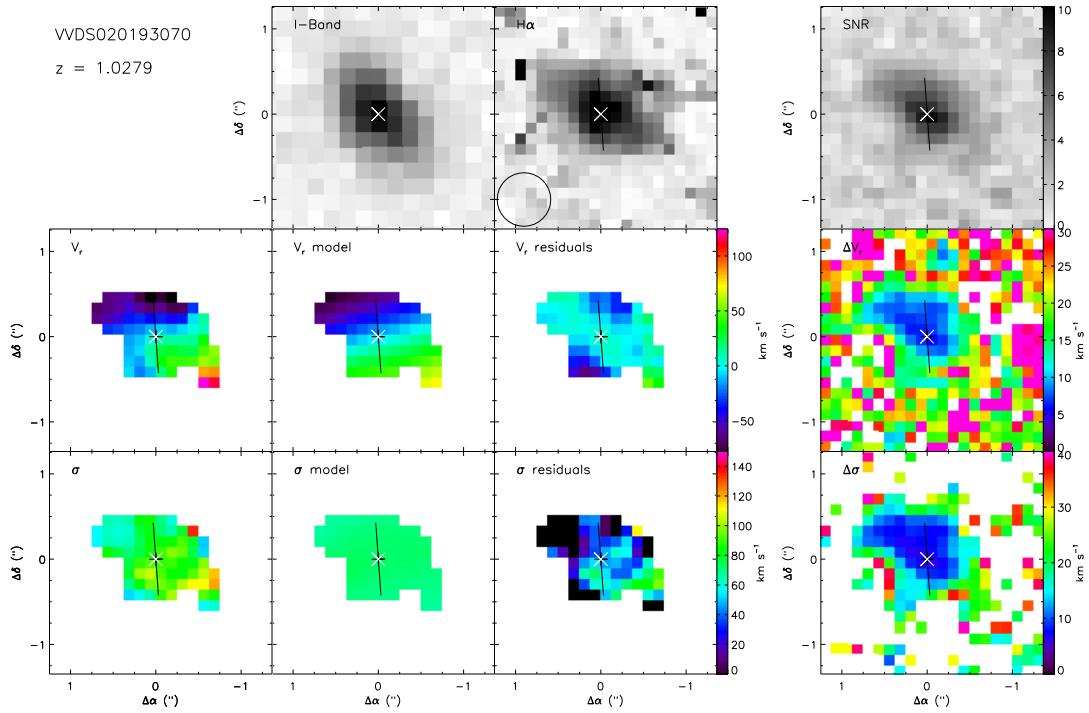


Fig. B.10. Maps for VVDS020193070. Same caption as Fig. B.1.

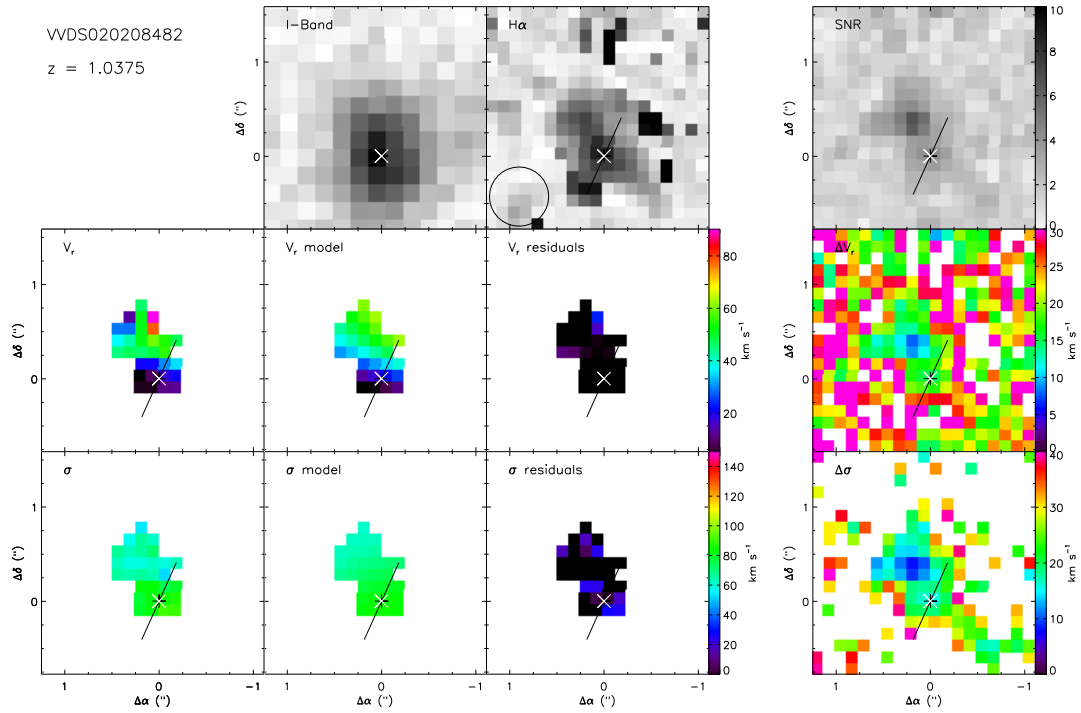


Fig. B.11. Maps for VVDS020208482. Same caption as Fig. B.1.

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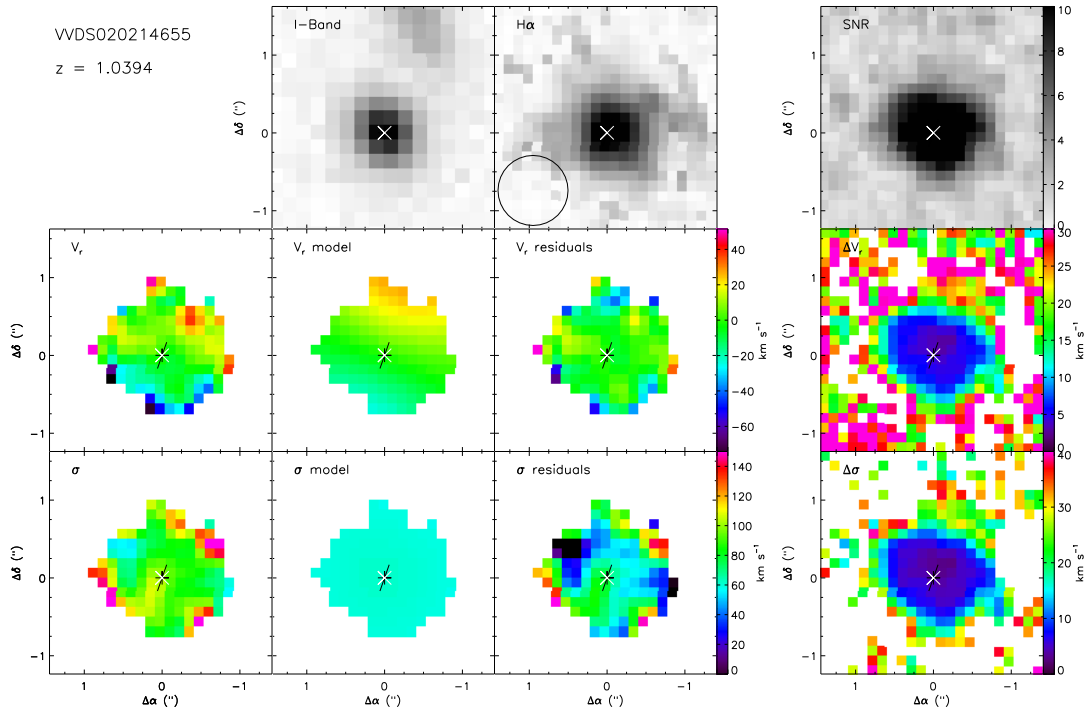


Fig. B.12. Maps for WVDS020214655. Same caption as Fig. B.1.

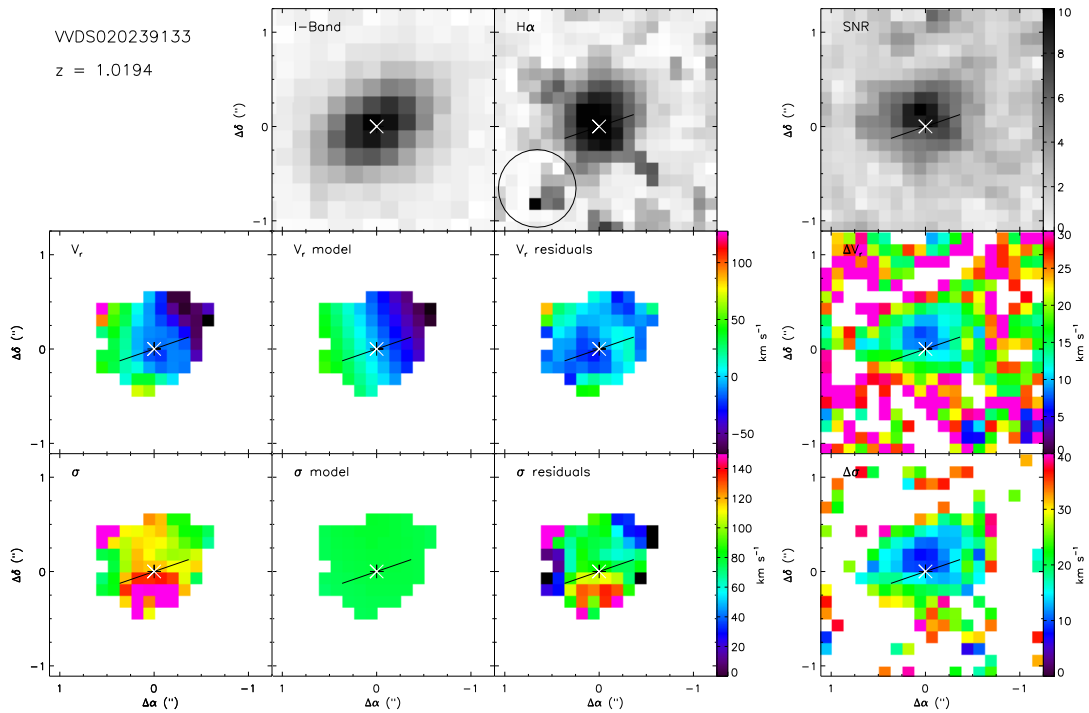


Fig. B.13. Maps for WVDS020239133. Same caption as Fig. B.1.

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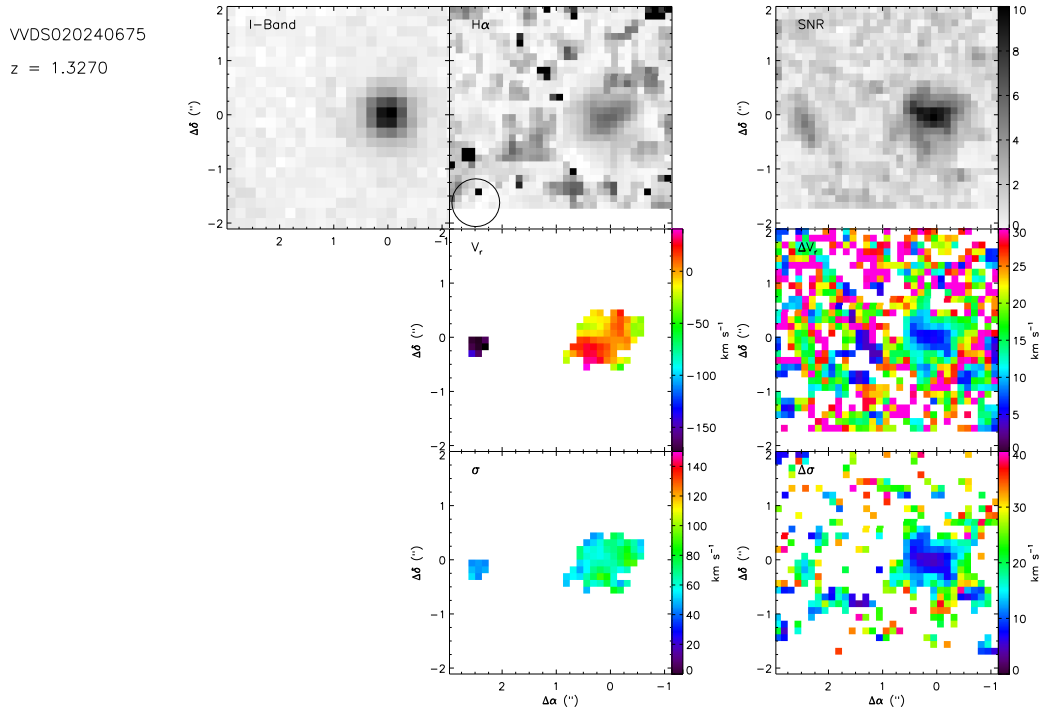


Fig. B.14. Maps for WVDS020240675 system. Same caption as Fig. B.2.

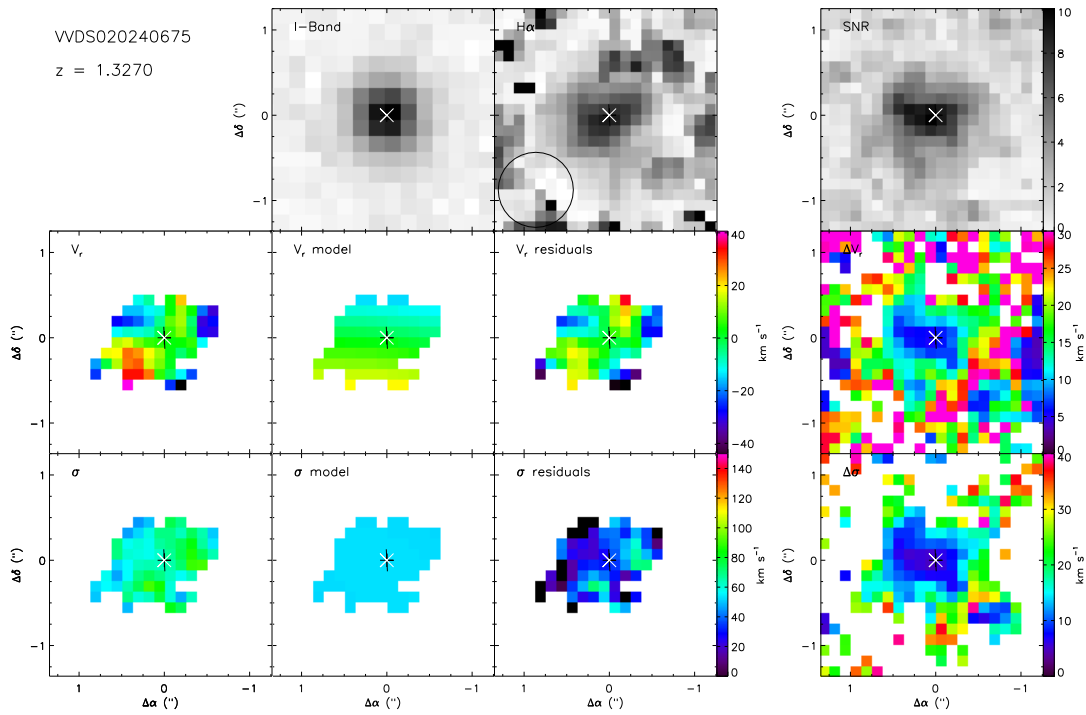


Fig. B.15. Maps for WVDS020240675. Same caption as Fig. B.1.

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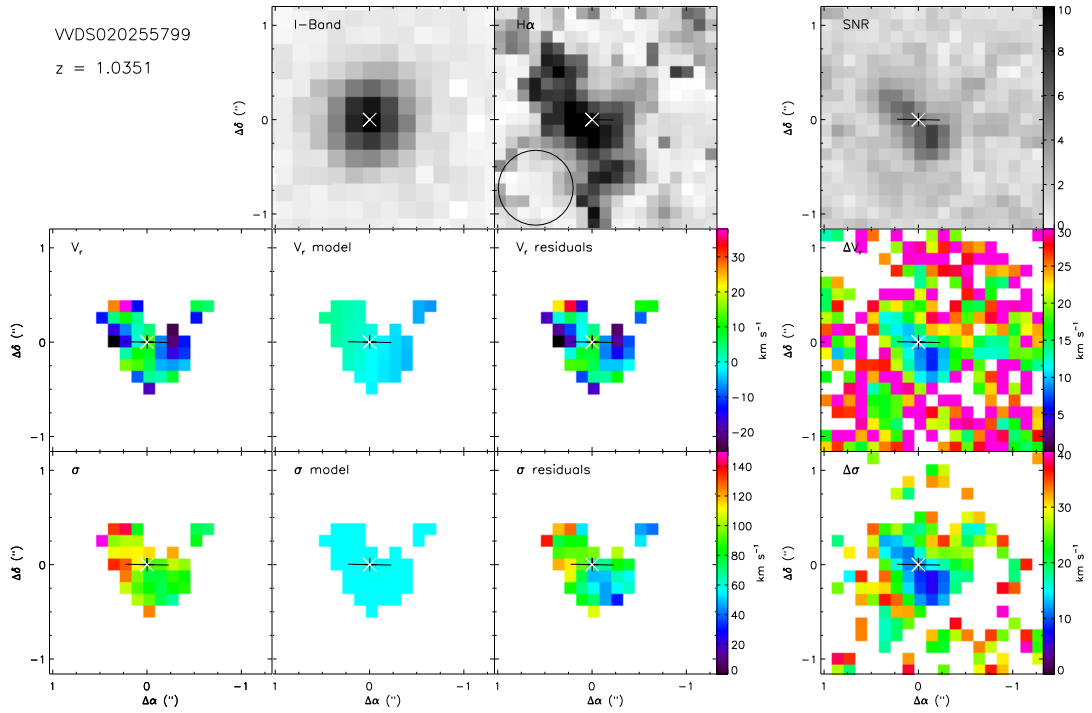


Fig. B.16. Maps for WVDS020255799. Same caption as Fig. B.1.

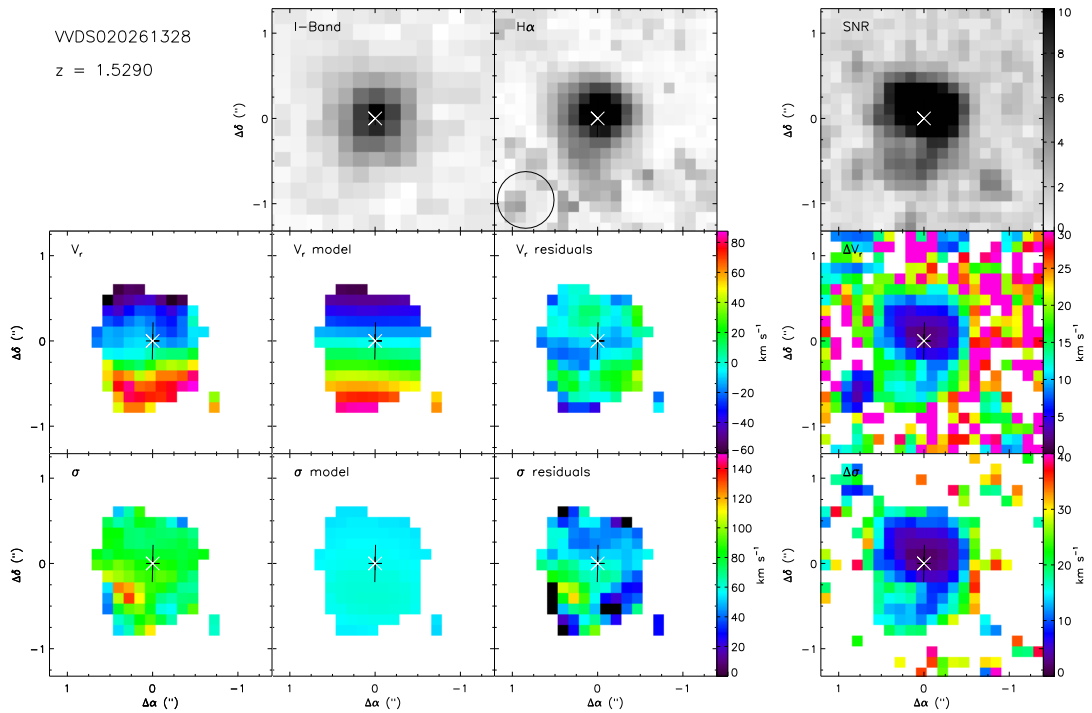


Fig. B.17. Maps for WVDS020261328. Same caption as Fig. B.1.

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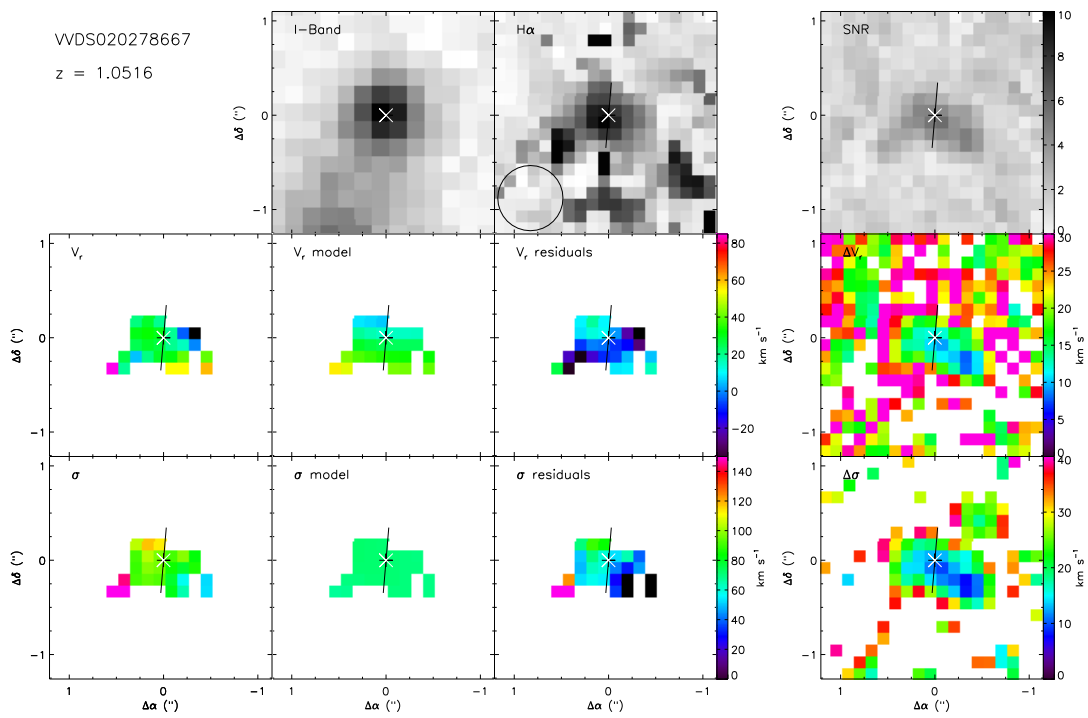


Fig. B.18. Maps for VVDS020278667. Same caption as Fig. B.1.

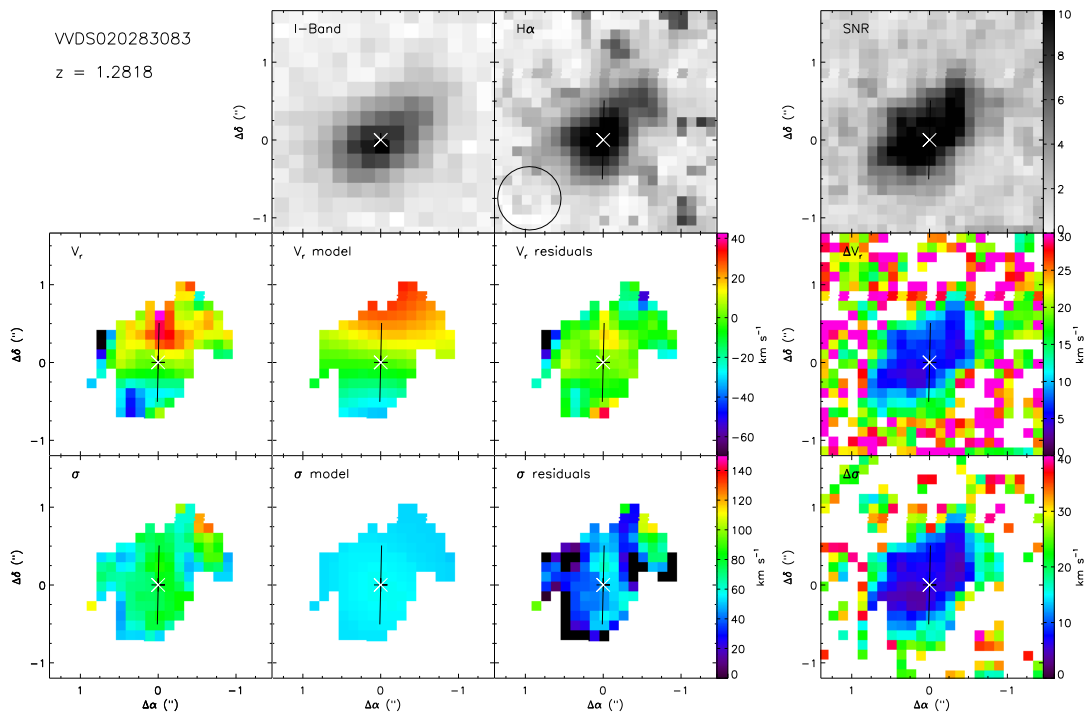


Fig. B.19. Maps for VVDS020283083. Same caption as Fig. B.1.

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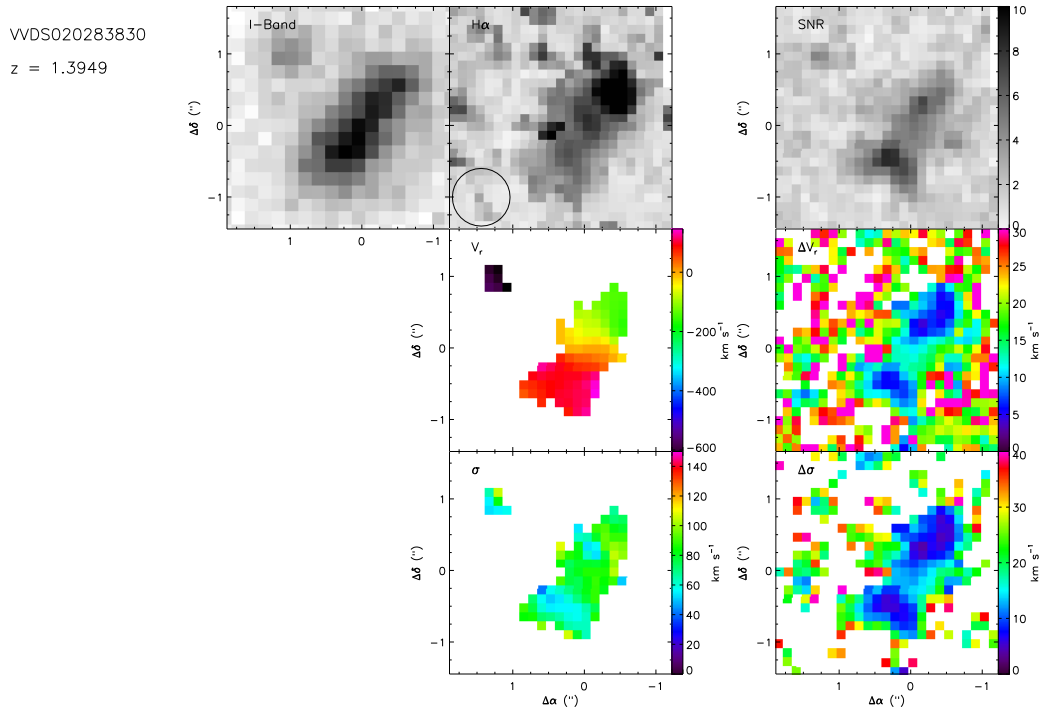


Fig. B.20. Maps for VVDS020283830 system. Same caption as Fig. B.2.

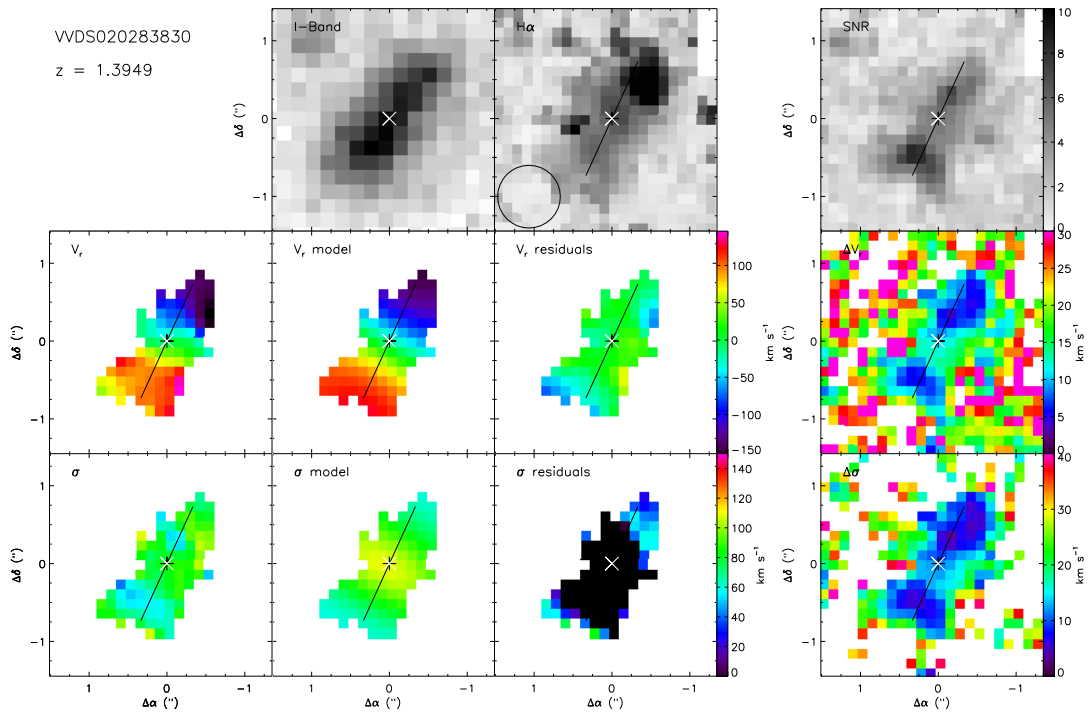


Fig. B.21. Maps for VVDS020283830. Same caption as Fig. B.1.

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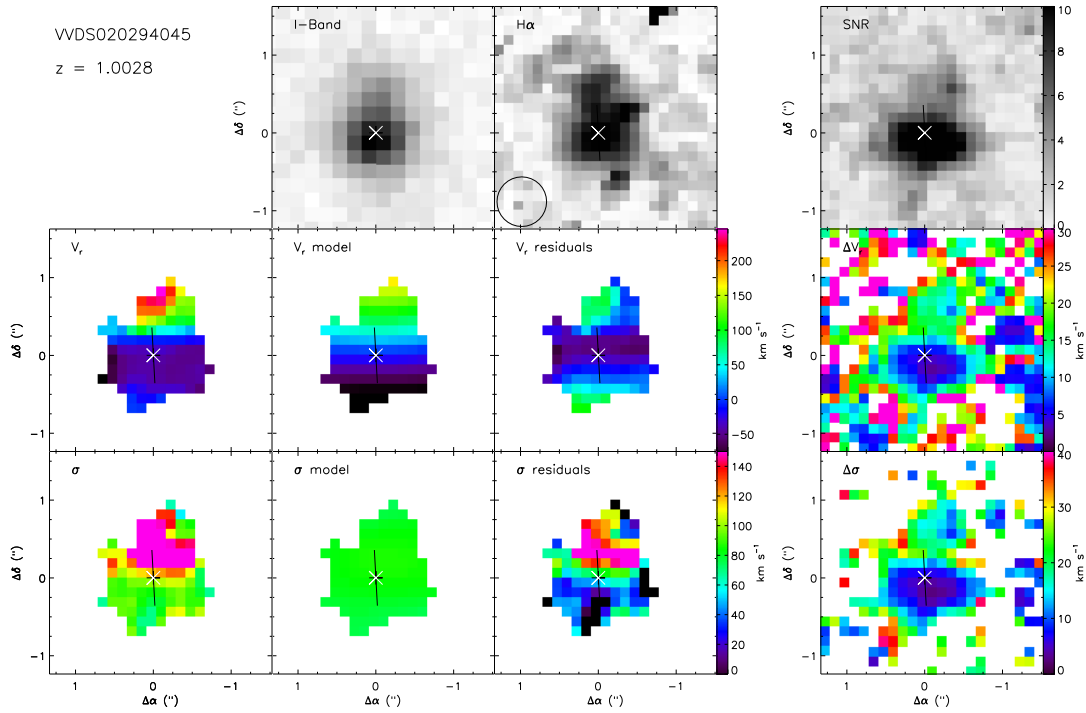


Fig. B.22. Maps for VVDS020294045. Same caption as Fig. B.1.

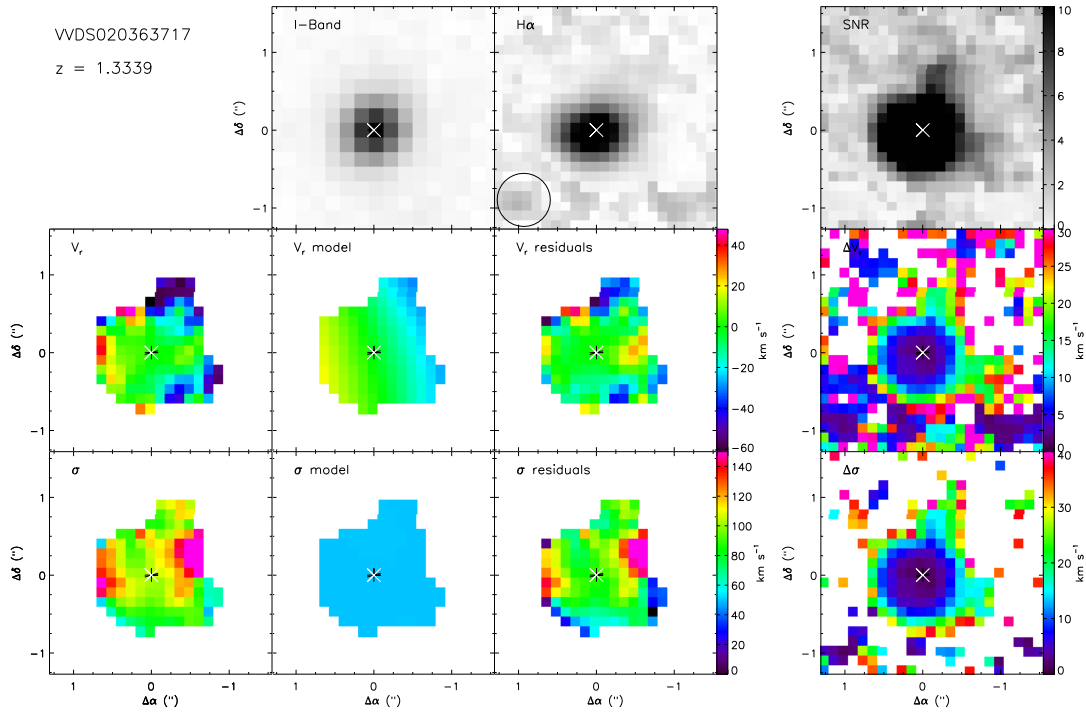


Fig. B.23. Maps for VVDS020363717. Same caption as Fig. B.1.

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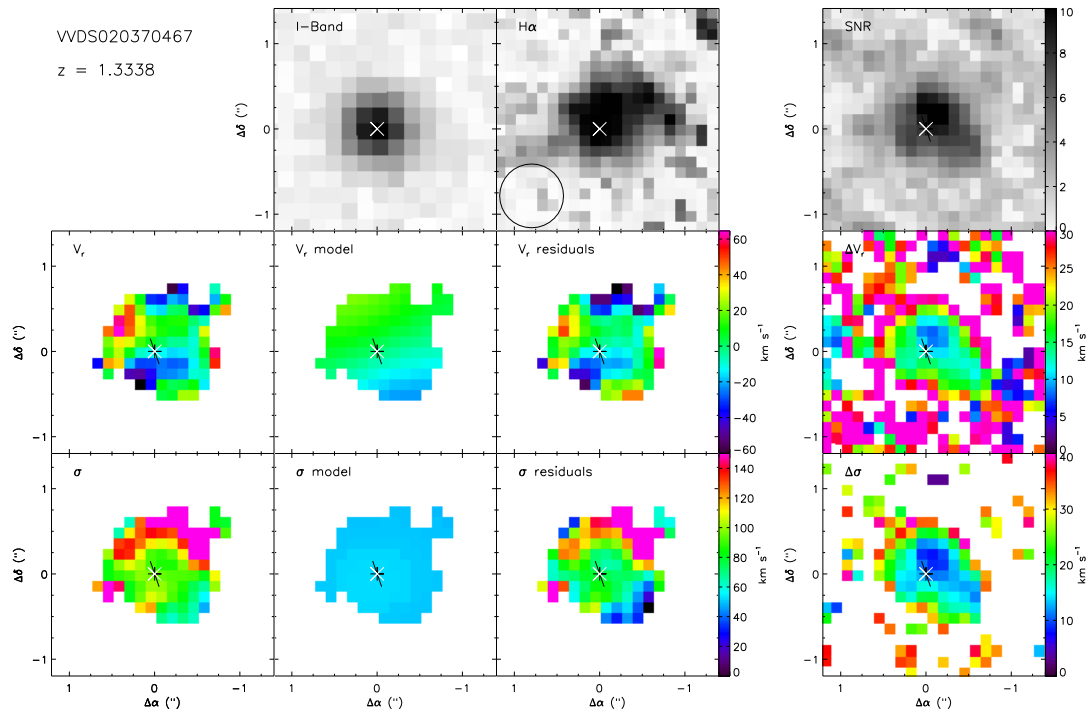


Fig. B.24. Maps for WVDS020370467. Same caption as Fig. B.1.

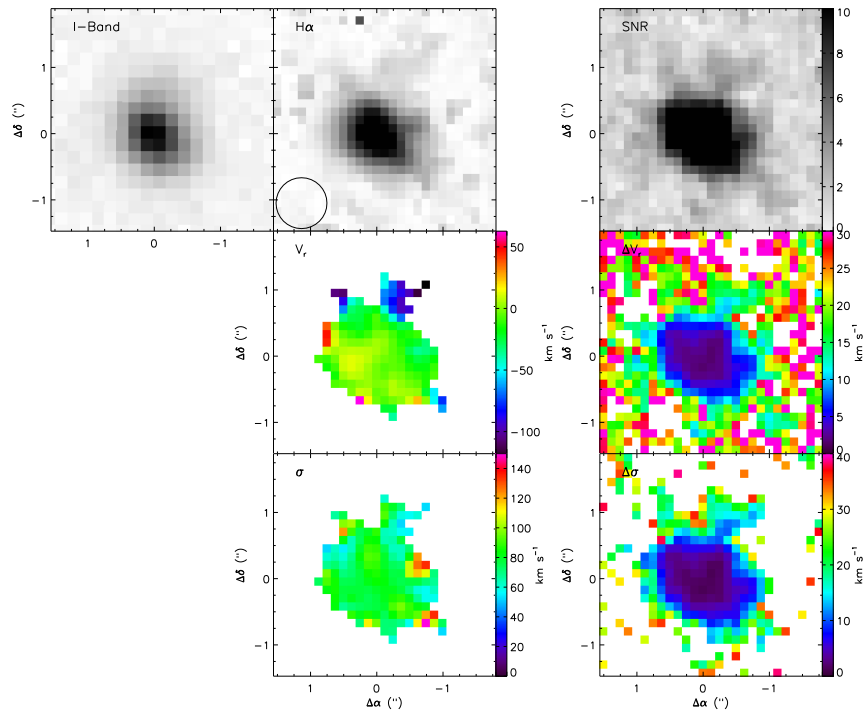


Fig. B.25. Maps for WVDS020386743 system. Same caption as Fig. B.2.

A&A 539, A92 (2012)

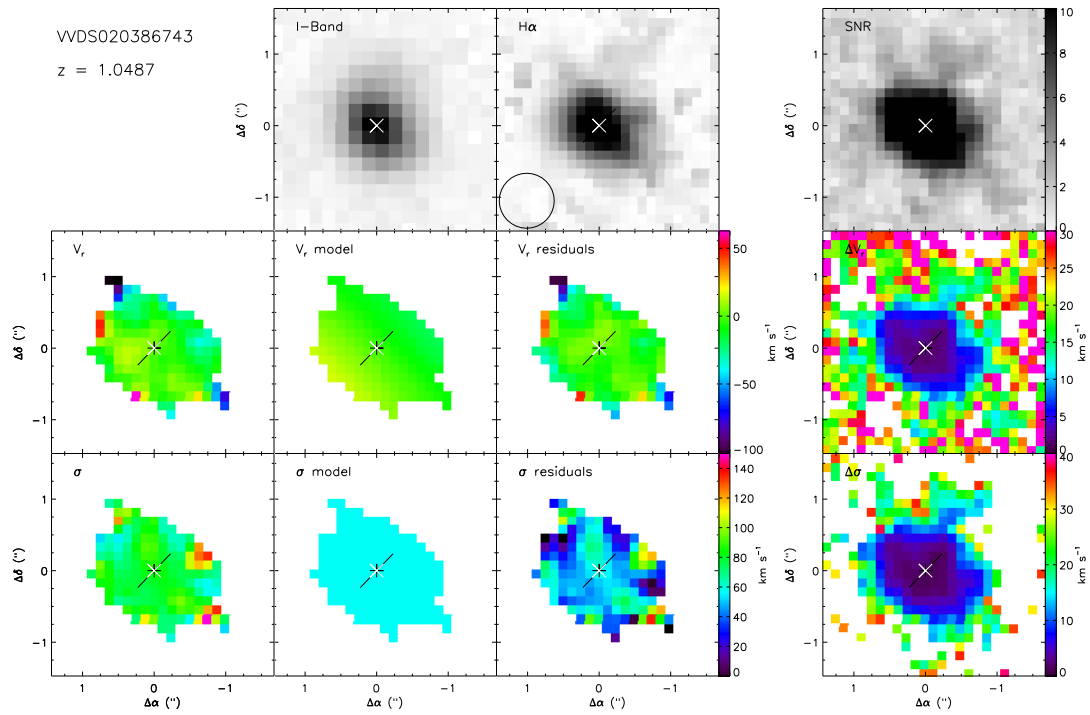


Fig. B.26. Maps for VVDS020386743. Same caption as Fig. B.1.

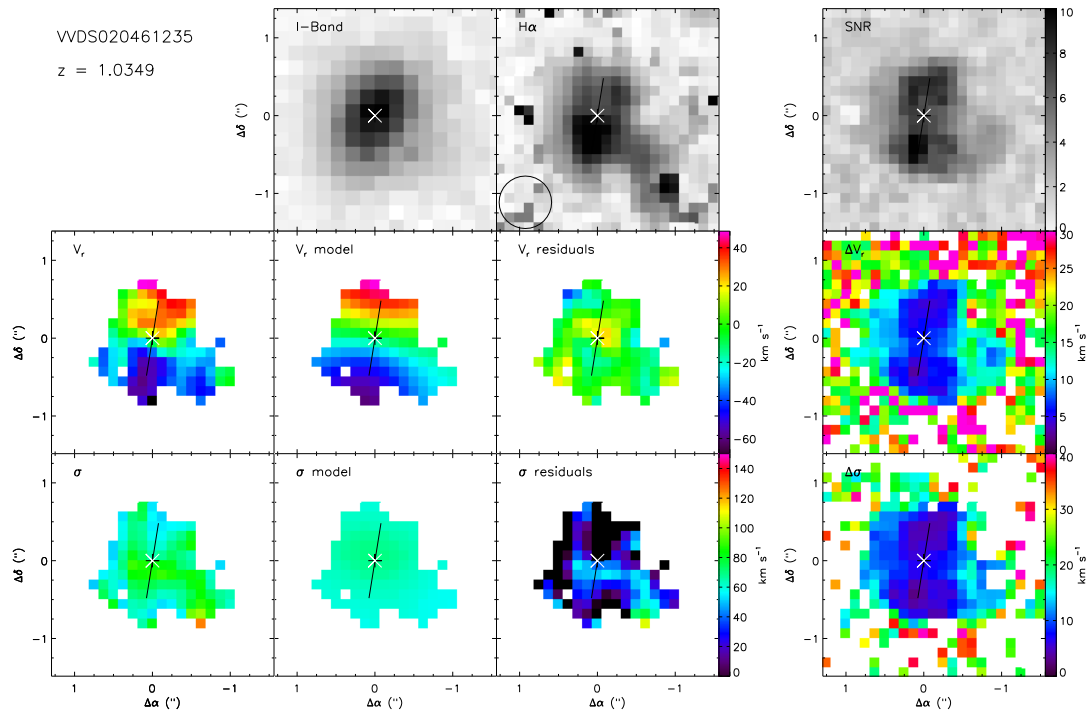


Fig. B.27. Maps for VVDS020461235. Same caption as Fig. B.1.

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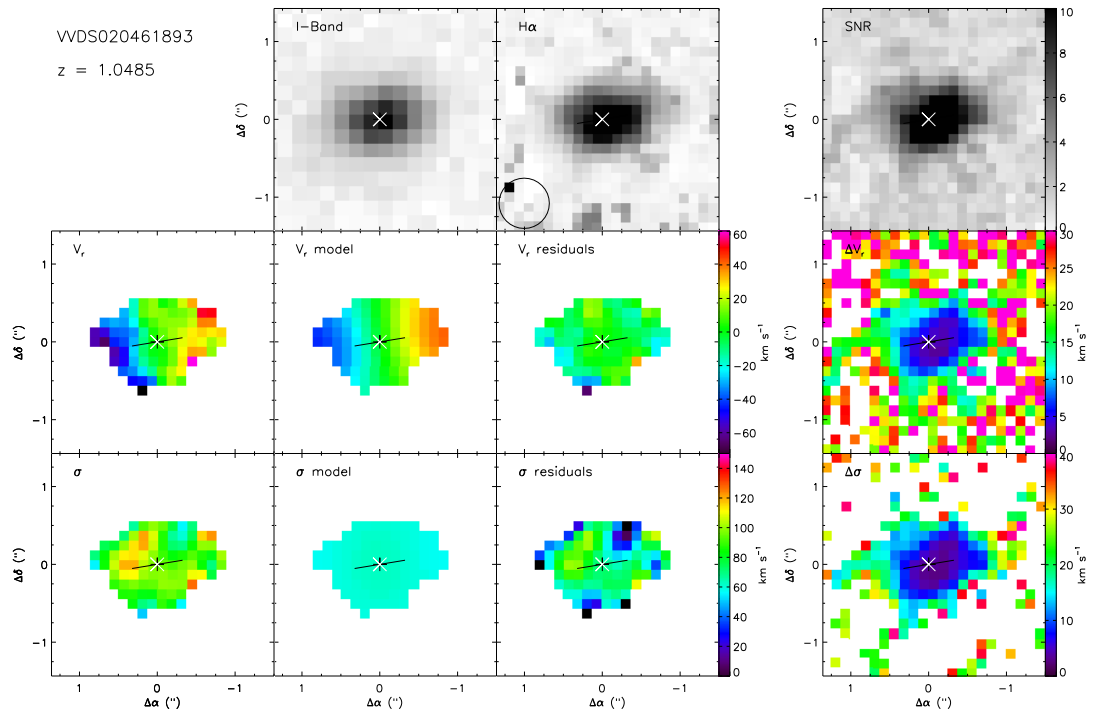


Fig. B.28. Maps for WVDS020461893. Same caption as Fig. B.1.

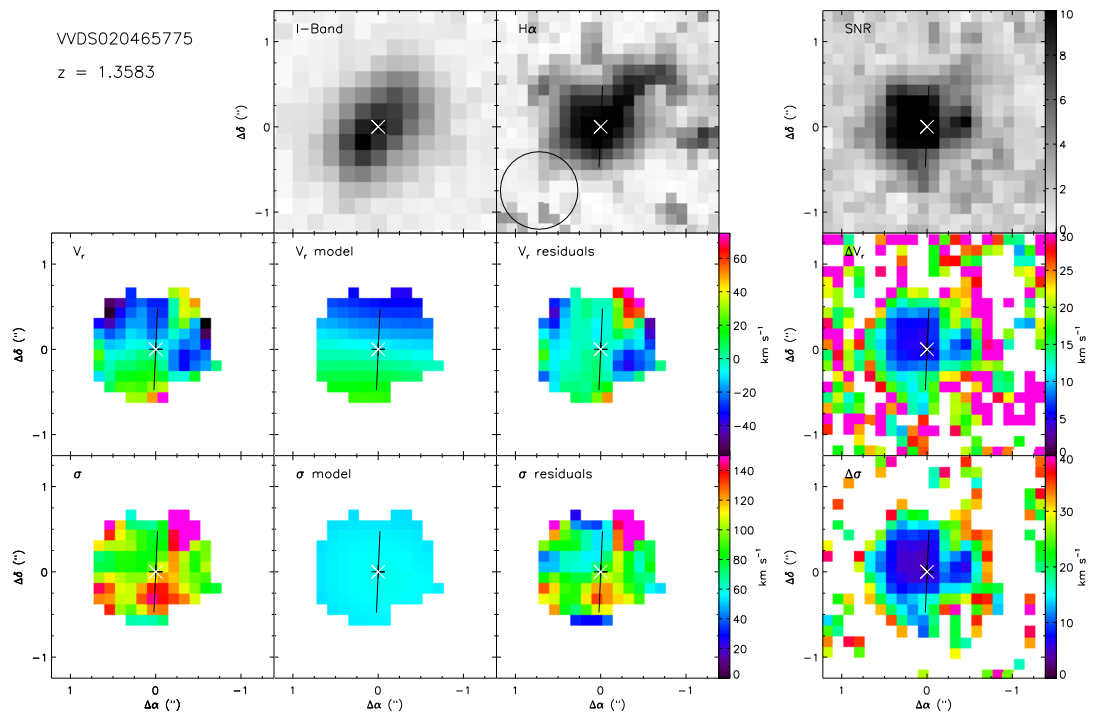


Fig. B.29. Maps for WVDS020465775. Same caption as Fig. B.1.

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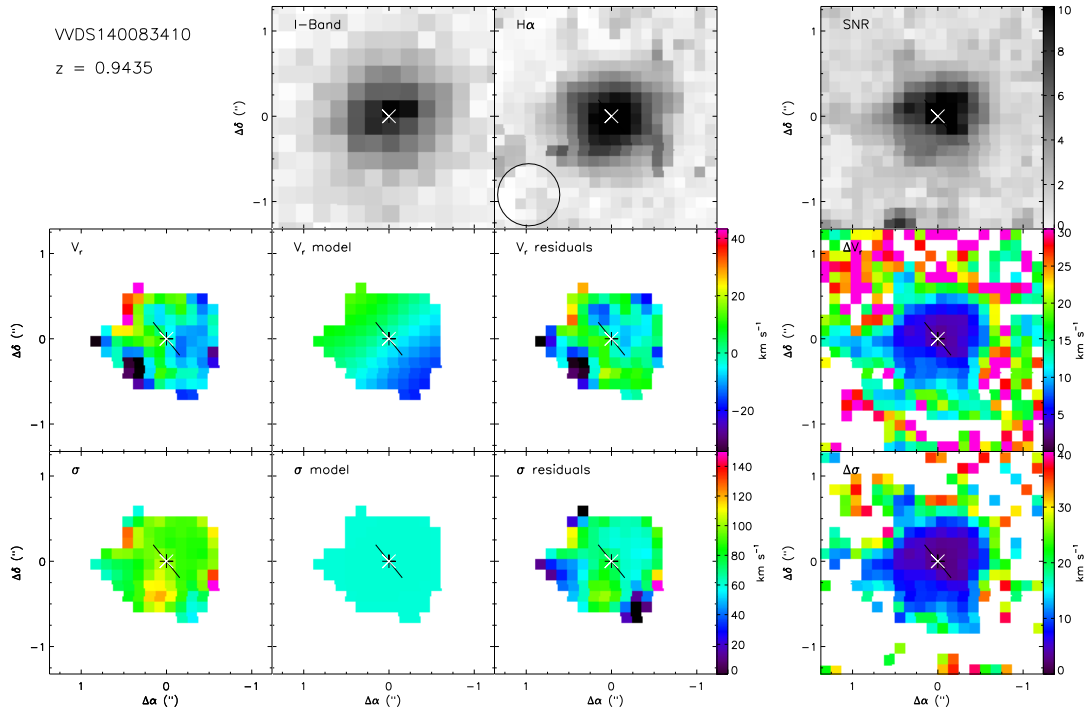


Fig. B.30. Maps for VVDS140083410. Same caption as Fig. B.1.

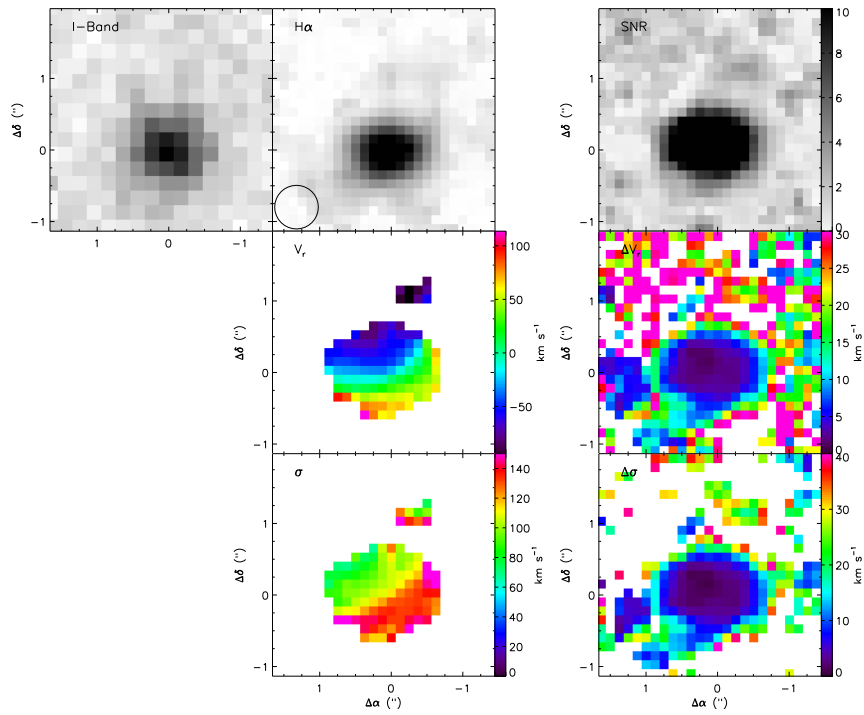


Fig. B.31. Maps for VVDS140096645 system. Same caption as Fig. B.2.

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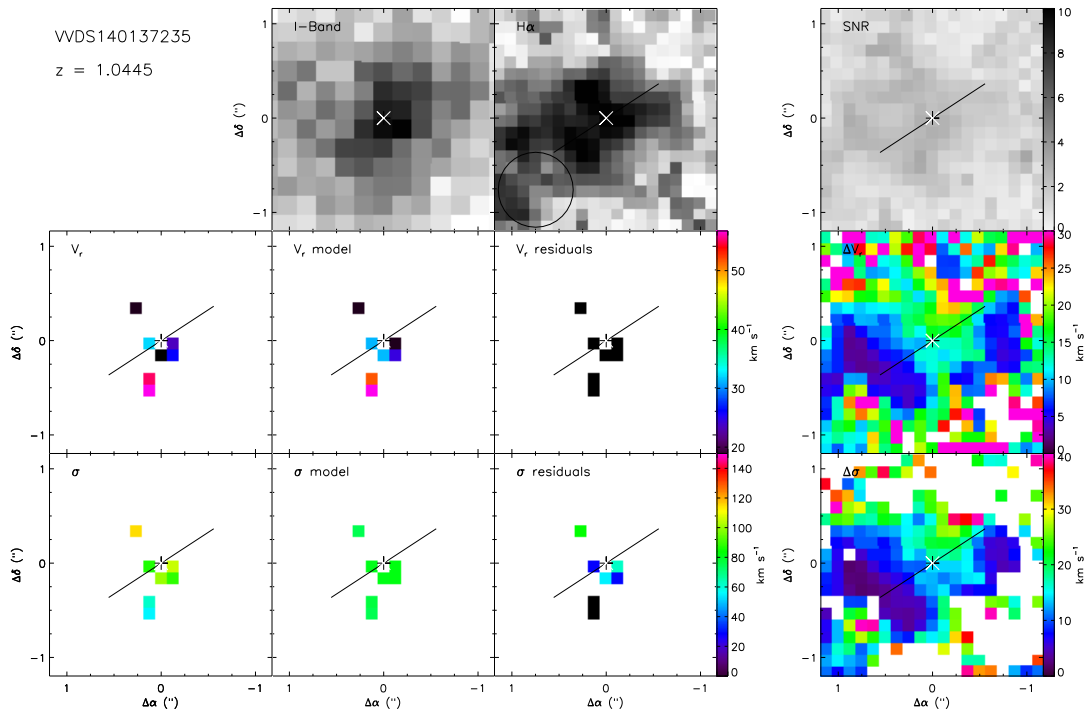


Fig. B.34. Maps for VVDS140137235. Same caption as Fig. B.1.

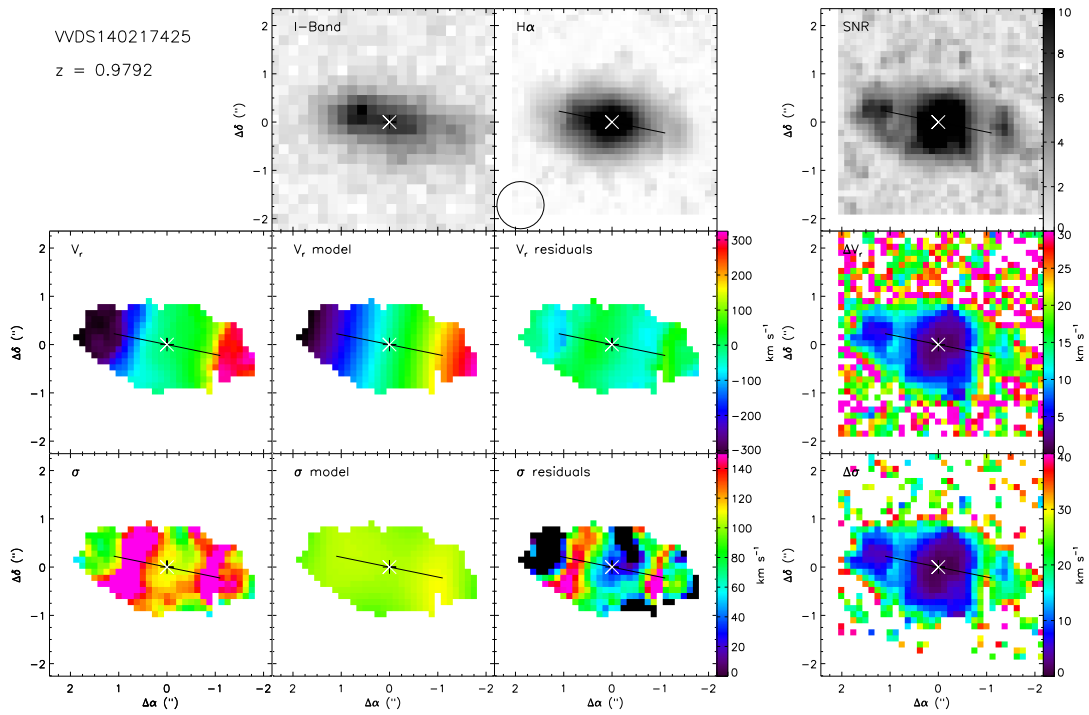


Fig. B.35. Maps for VVDS140217425. Same caption as Fig. B.1.

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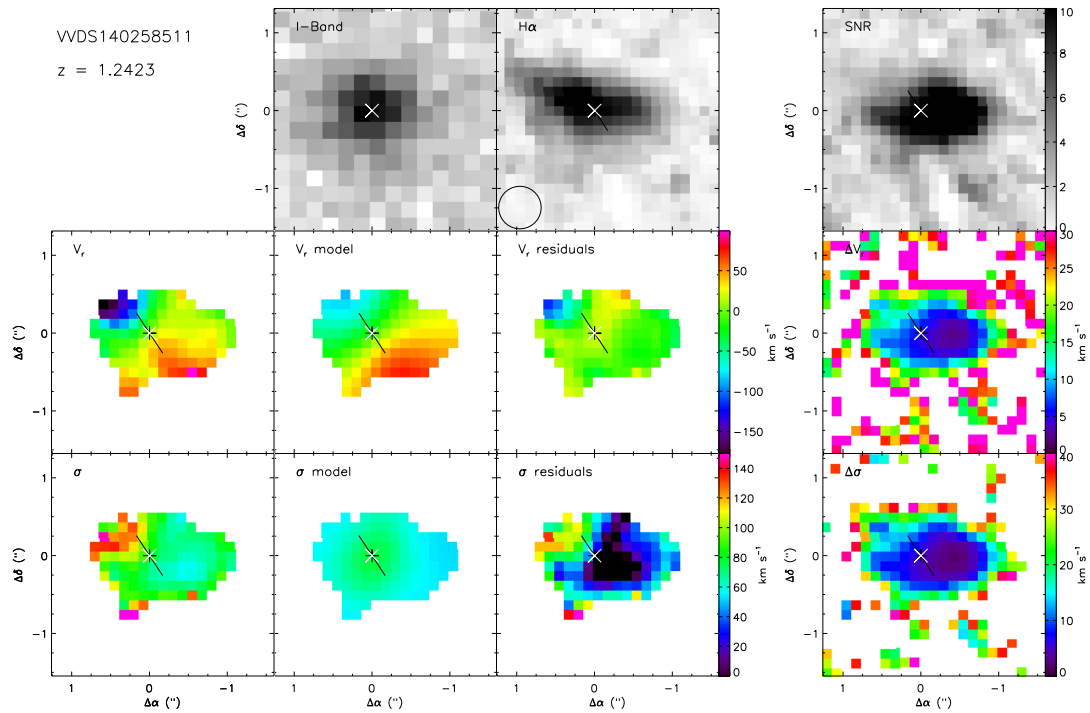


Fig. B.36. Maps for VVDS140258511. Same caption as Fig. B.1.

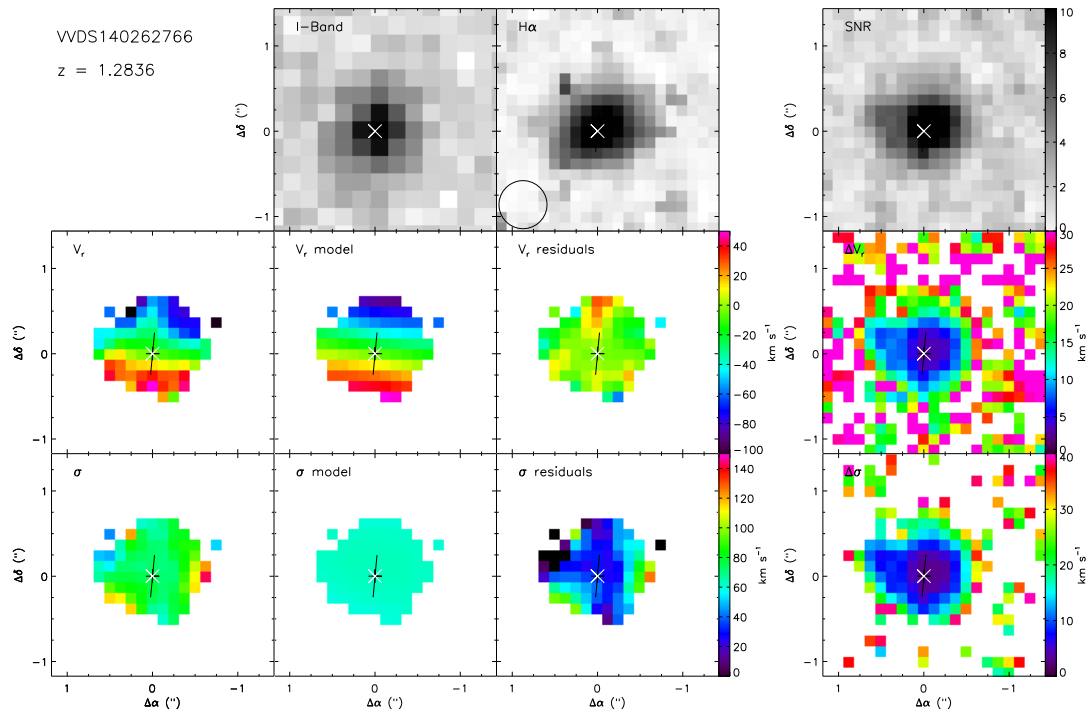


Fig. B.37. Maps for VVDS140262766. Same caption as Fig. B.1.

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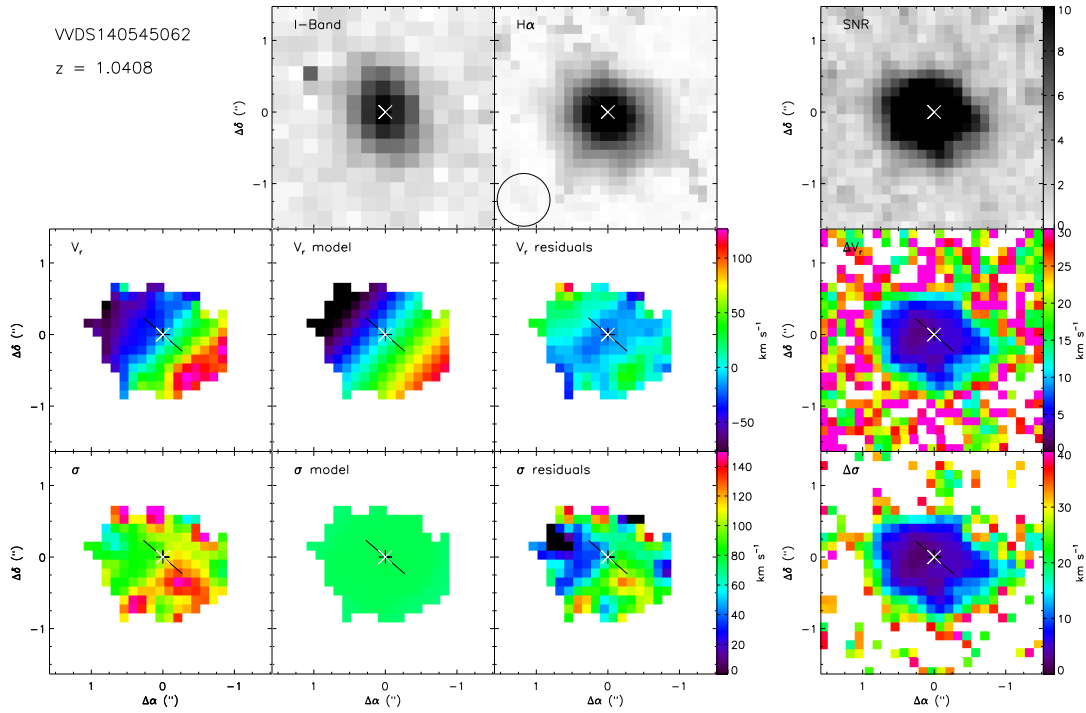


Fig. B.38. Maps for VVDS140545062. Same caption as Fig. B.1.

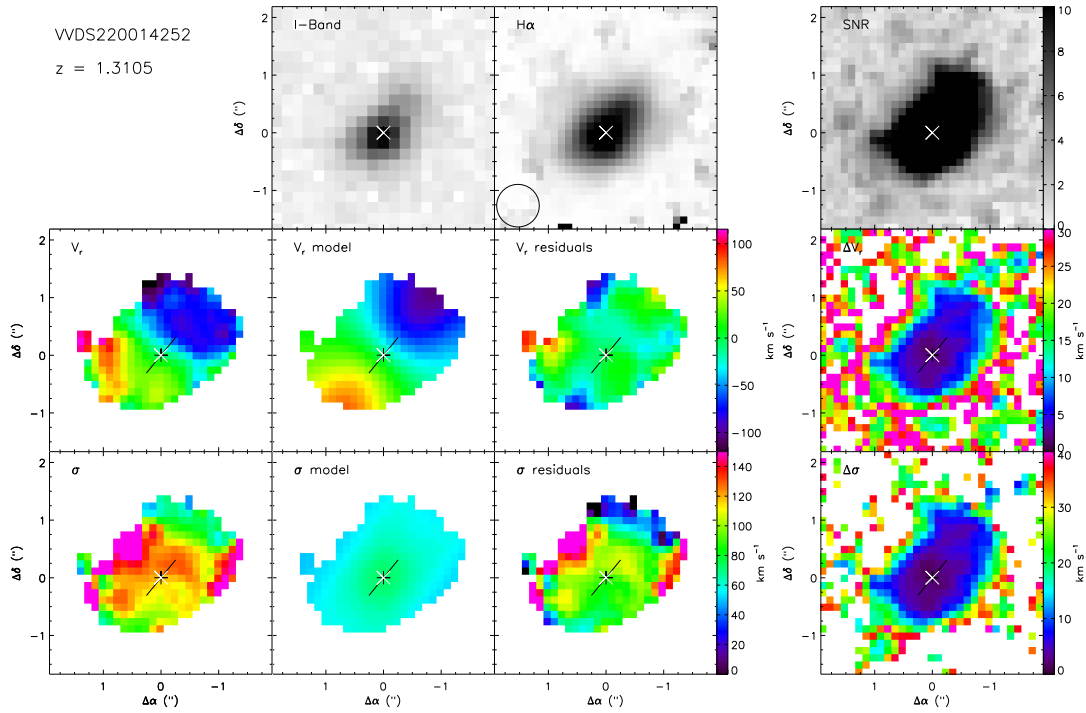


Fig. B.39. Maps for VVDS220014252. Same caption as Fig. B.1.

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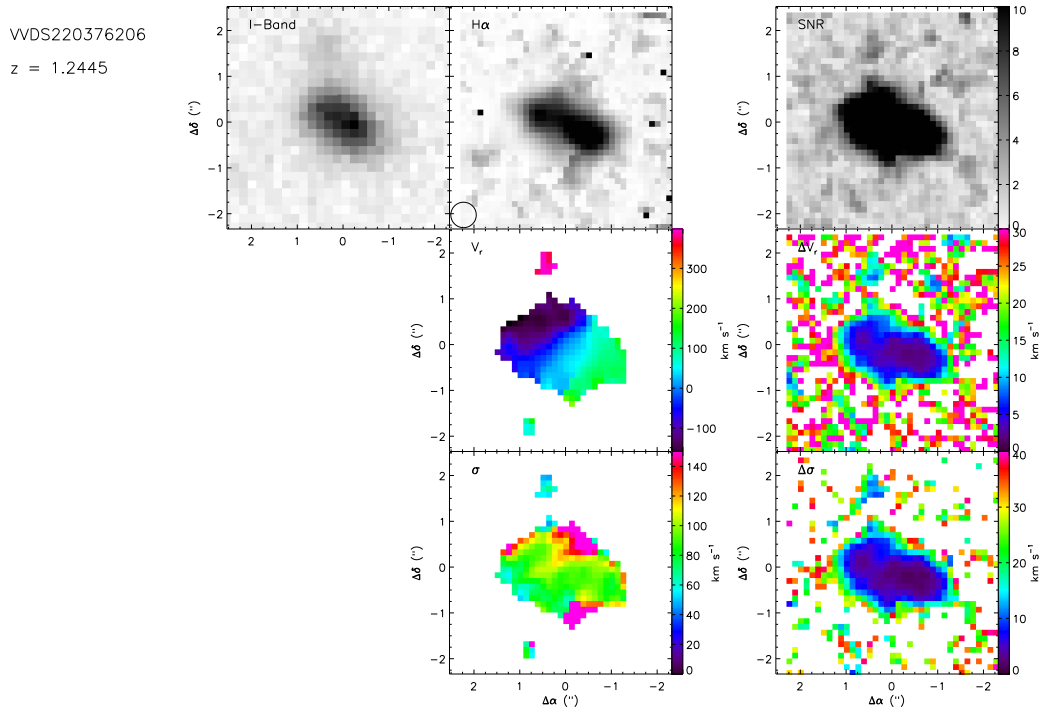


Fig. B.42. Maps for VVDS220376206 system. Same caption as Fig. B.2.

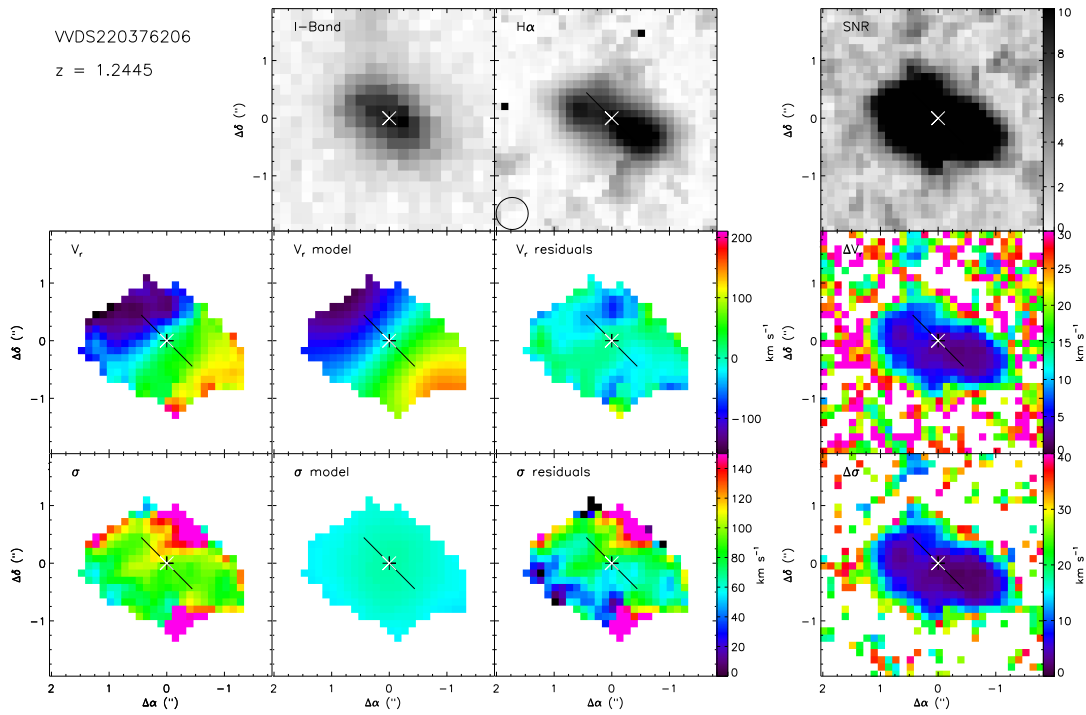


Fig. B.43. Maps for VVDS220376206. Same caption as Fig. B.1.

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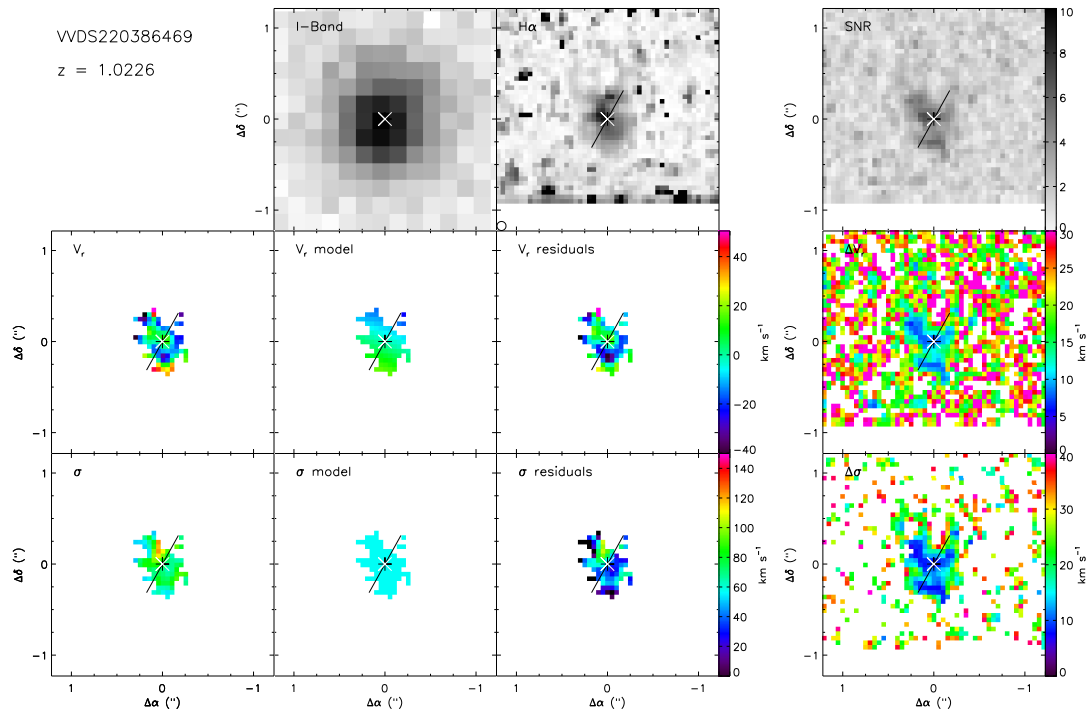


Fig. B.44. Maps for WDS220386469. Same caption as Fig. B.1.

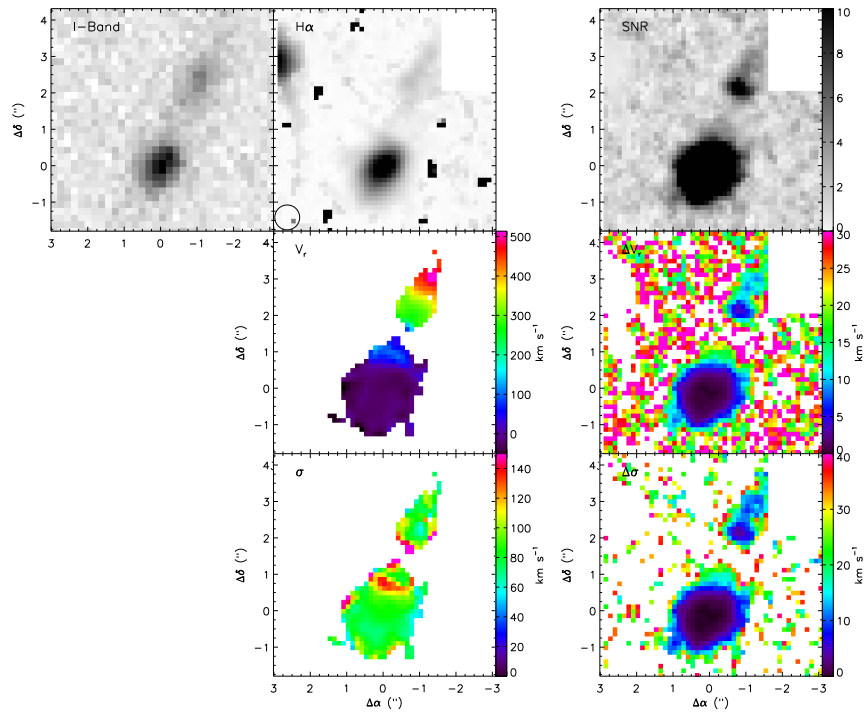


Fig. B.45. Maps for WDS220397579 system. Same caption as Fig. B.2.

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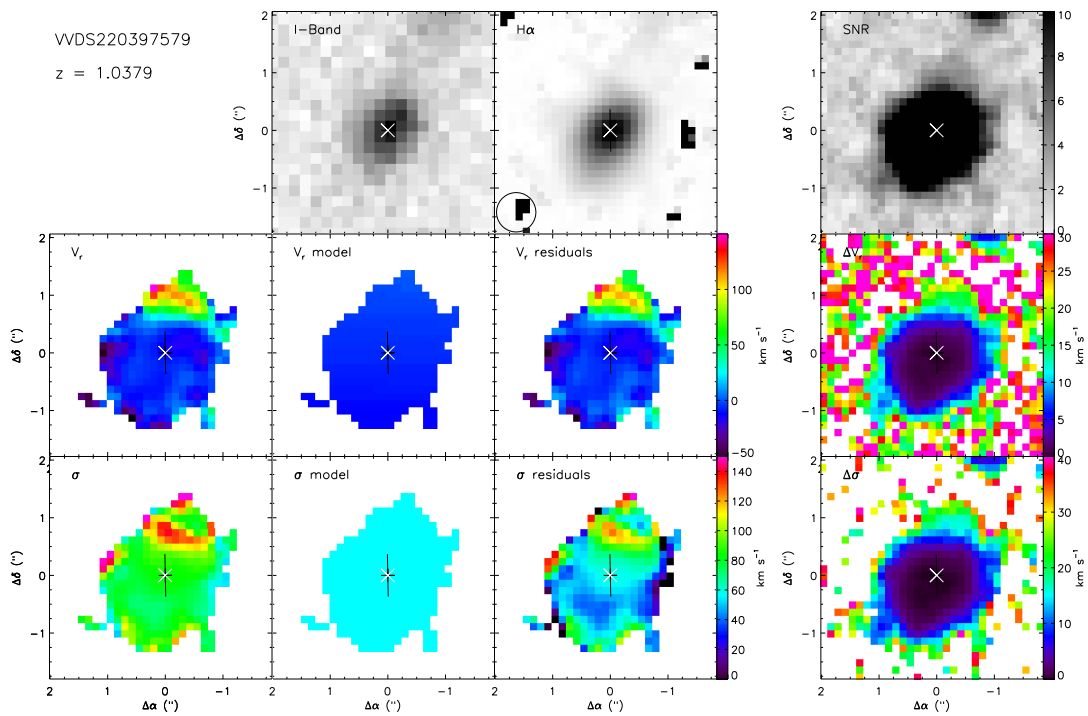


Fig. B.46. Maps for VVDS220397579. Same caption as Fig. B.1.

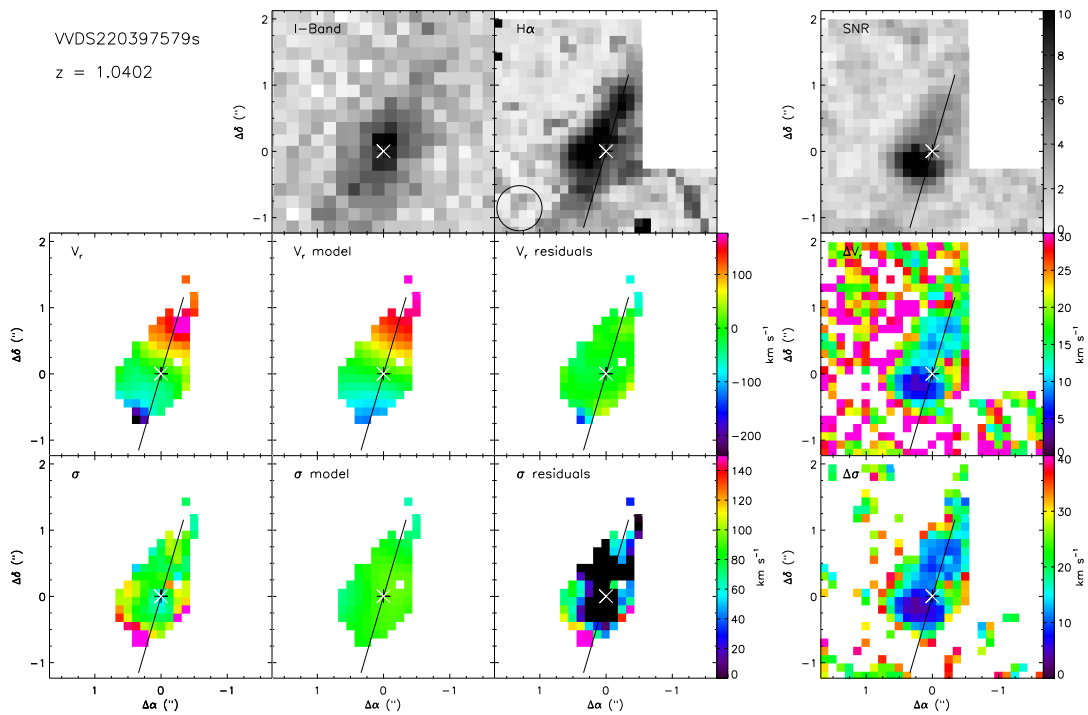


Fig. B.47. Maps for VVDS220397579s companion. Same caption as Fig. B.1.

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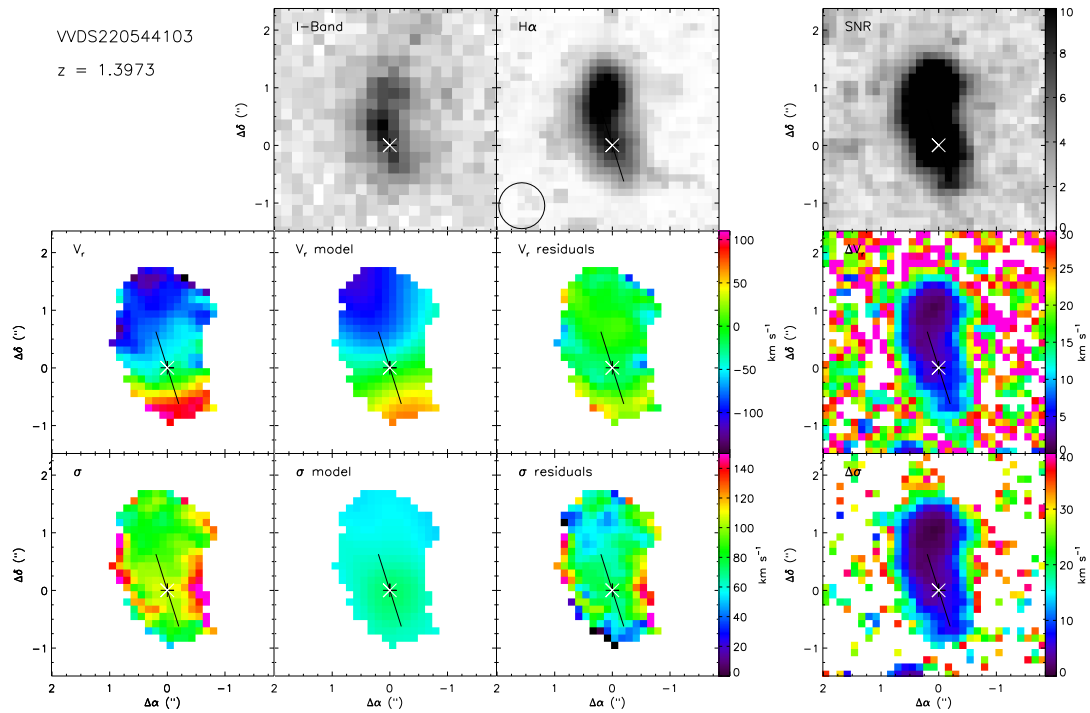


Fig. B.48. Maps for VVDS220544103. Same caption as Fig. B.1.

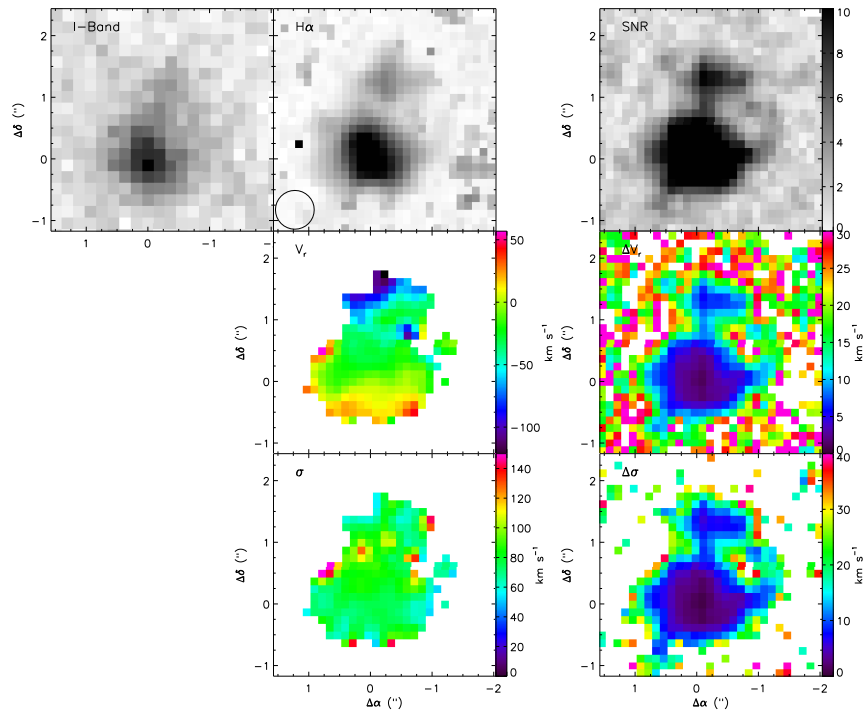


Fig. B.49. Maps for VVDS220544394 system. Same caption as Fig. B.2.

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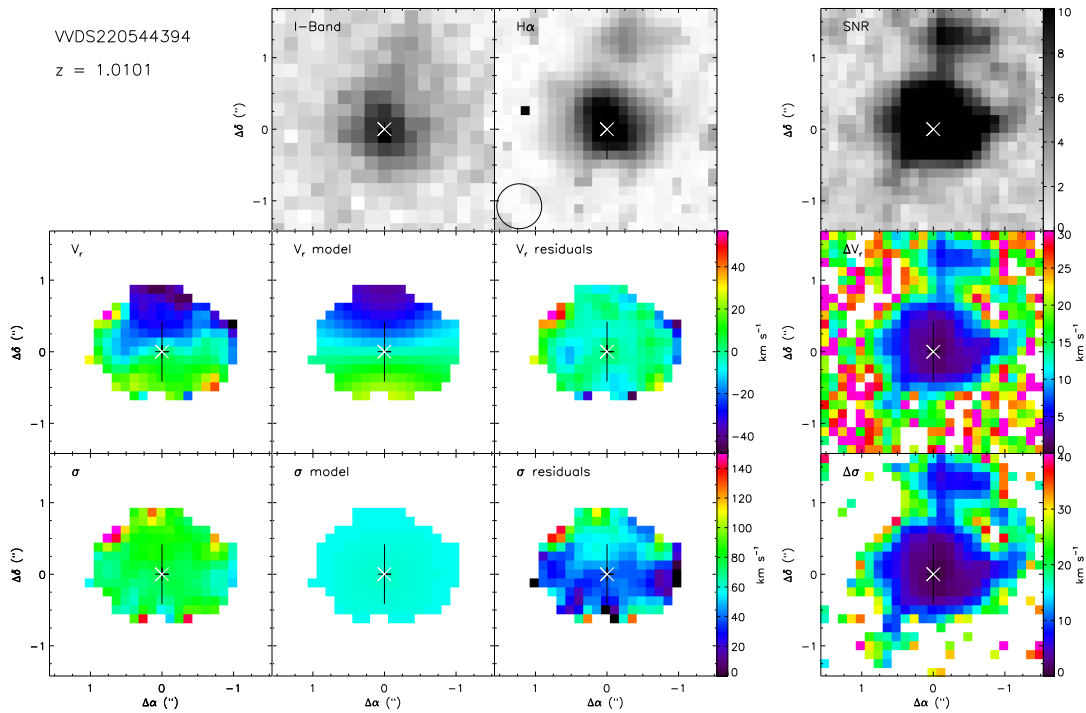


Fig. B.50. Maps for VVDS220544394. Same caption as Fig. B.1.

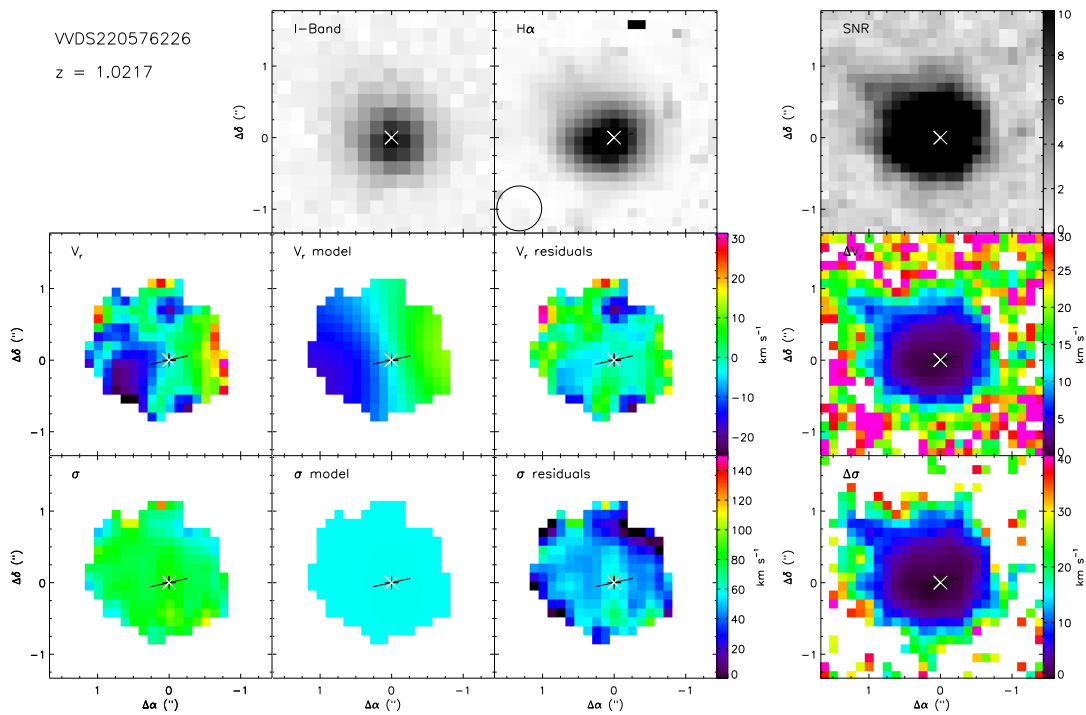


Fig. B.51. Maps for VVDS220576226. Same caption as Fig. B.1.

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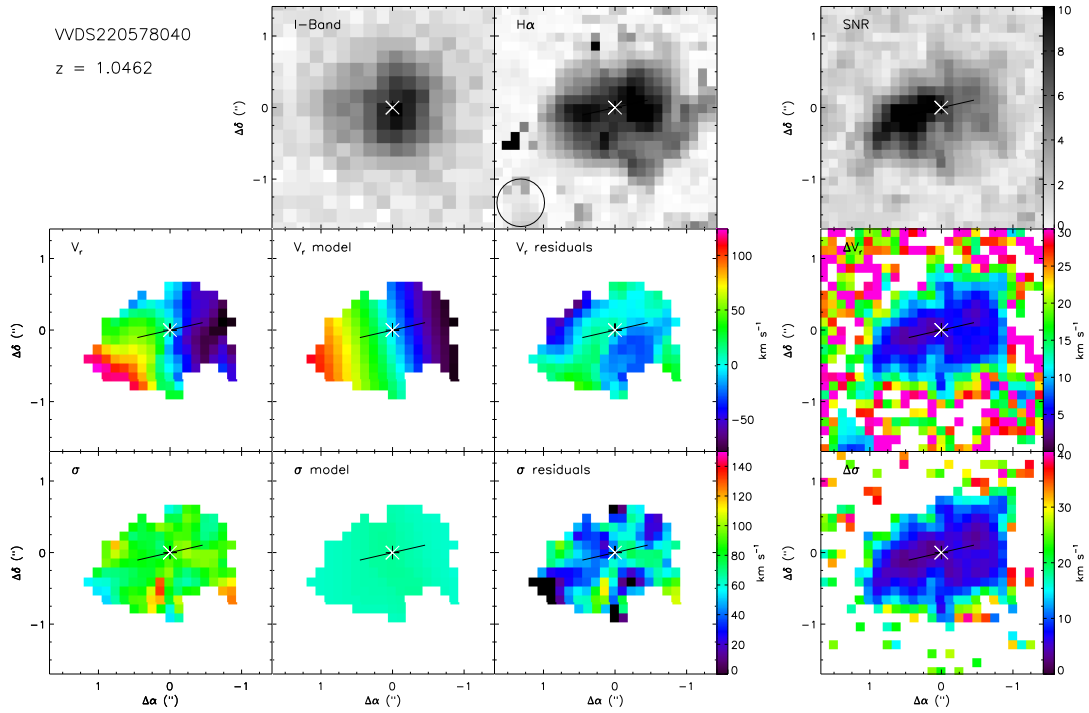


Fig. B.52. Maps for WVDS220578040. Same caption as Fig. B.1.

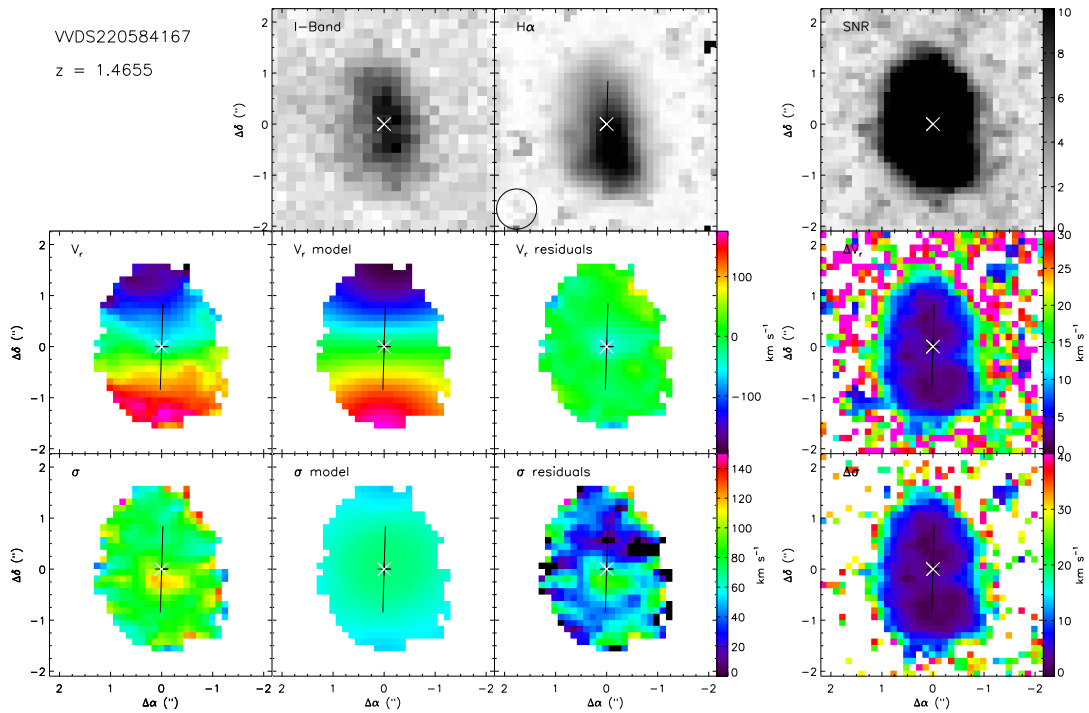


Fig. B.53. Maps for WVDS220584167. Same caption as Fig. B.1.

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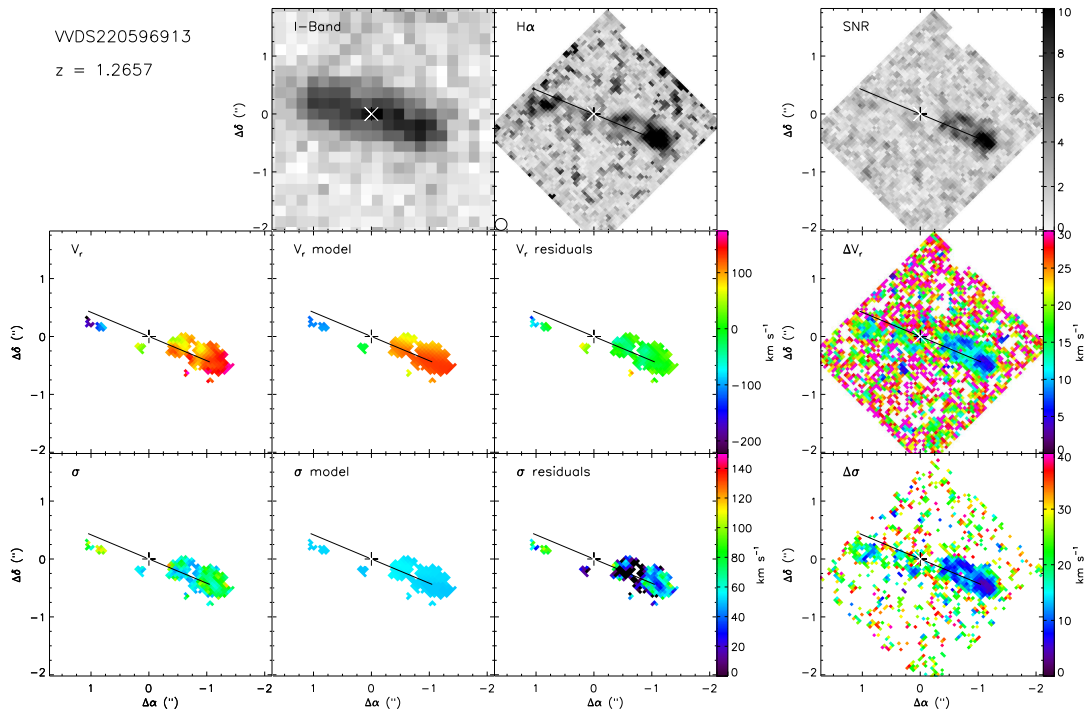


Fig. B.54. Maps for VVDS220596913. Same caption as Fig. B.1.

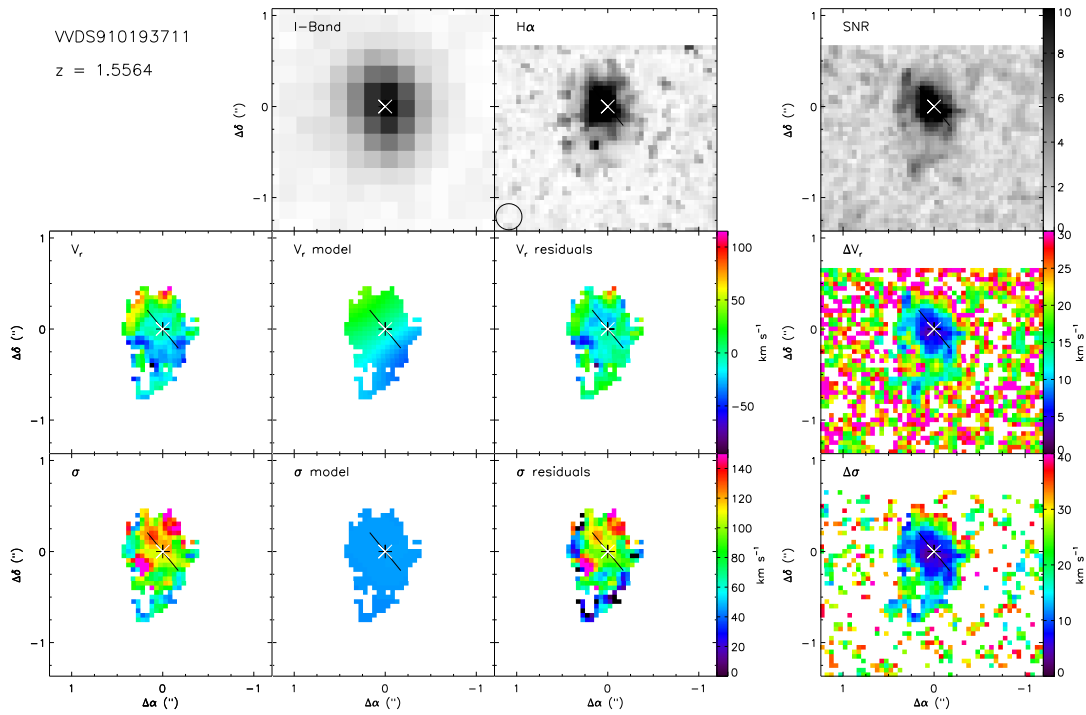


Fig. B.55. Maps for VVDS910193711. Same caption as Fig. B.1.

B. Epinat et al.: MASSIV: Mass Assembly Survey with SINFONI in VVDS. II.

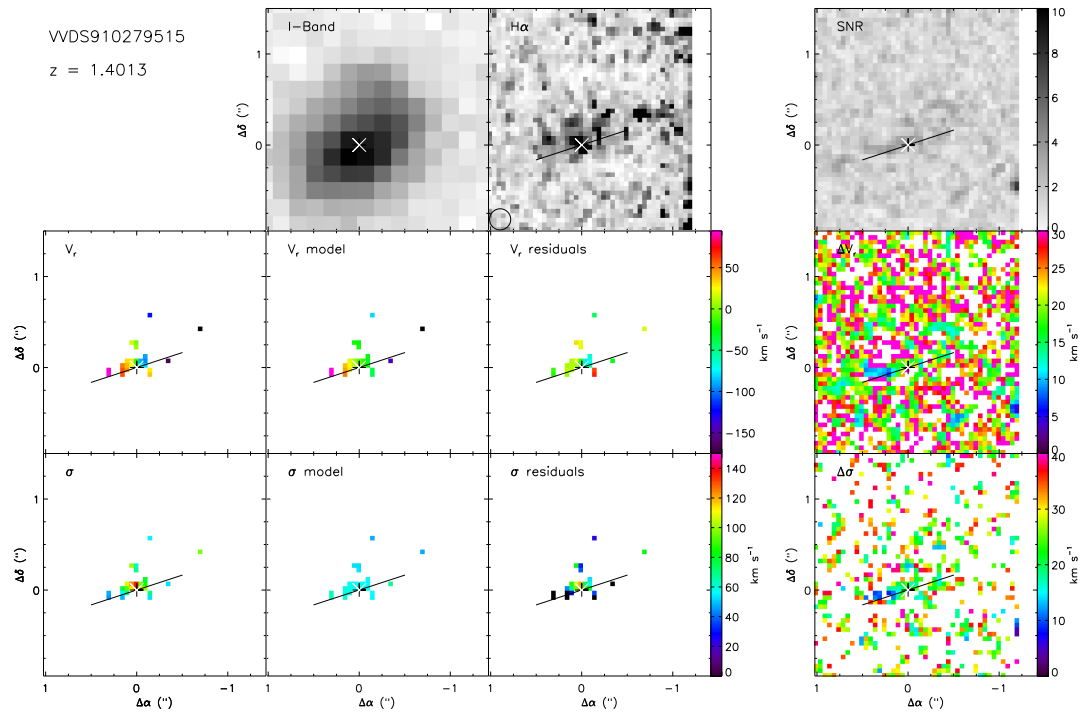


Fig. B.56. Maps for VVDS910279515. Same caption as Fig. B.1.

B.2 Article IV : MASSIV V. The major merger rate of star-forming galaxies at $0.9 < z < 1.8$ from IFS-based close pairs

MASSIV: Mass Assembly Survey with SINFONI in VVDS***V. The major merger rate of star-forming galaxies at $0.9 < z < 1.8$ from IFS-based close pairs**C. López-Sanjuan^{1,2}, O. Le Fèvre¹, L. A. M. Tasca¹, B. Epinat^{1,3,4}, P. Amram¹, T. Contini^{3,4}, B. Garilli⁵, M. Kissler-Patig⁶, J. Moutaka^{3,4}, L. Paioro⁵, V. Perret¹, J. Queyrel^{3,4}, L. Tresse¹, D. Vergani⁷, and C. Divoy^{3,4}¹ Aix Marseille Université, CNRS, LAM (Laboratoire d'Astrophysique de Marseille) UMR 7326, 13388, Marseille, France² Centro de Estudios de Física del Cosmos de Aragón, Plaza San Juan 1, planta 2, 44001, Teruel, Spain

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ABSTRACT

Context. The contribution of the merging process to the early phase of galaxy assembly at $z > 1$ and, in particular, to the build-up of the red sequence, still needs to be accurately assessed.

Aims. We aim to measure the major merger rate of star-forming galaxies at $0.9 < z < 1.8$, using close pairs identified from integral field spectroscopy (IFS).

Methods. We use the velocity field maps obtained with SINFONI/VLT on the MASSIV sample, selected from the star-forming population in the VVDS. We identify physical pairs of galaxies from the measurement of the relative velocity and the projected separation (r_p) of the galaxies in the pair. Using the well constrained selection function of the MASSIV sample, we derive at a mean redshift up to $z = 1.54$ the gas-rich major merger fraction (luminosity ratio $\mu = L_2/L_1 \geq 1/4$), and the gas-rich major merger rate using merger time scales from cosmological simulations.

Results. We find a high gas-rich major merger fraction of $20.8^{+15.2}_{-6.8}\%$, $20.1^{+8.0}_{-5.1}\%$, and $22.0^{+13.7}_{-7.3}\%$ for close pairs with $r_p \leq 20h^{-1}$ kpc in redshift ranges $z = [0.94, 1.06]$, $[1.2, 1.5]$, and $[1.5, 1.8]$, respectively. This translates into a gas-rich major merger rate of $0.116^{+0.084}_{-0.038} \text{ Gyr}^{-1}$, $0.147^{+0.058}_{-0.037} \text{ Gyr}^{-1}$, and $0.127^{+0.079}_{-0.042} \text{ Gyr}^{-1}$ at $z = 1.03, 1.32$, and 1.54 , respectively. Combining our results with previous studies at $z < 1$, the gas-rich major merger rate evolves as $(1+z)^n$, with $n = 3.95 \pm 0.12$, up to $z = 1.5$. From these results we infer that $\sim 35\%$ of the star-forming galaxies with stellar masses $\overline{M}_* = 10^{10} - 10^{10.5} M_\odot$ have undergone a major merger since $z \sim 1.5$. We develop a simple model that shows that, assuming that all gas-rich major mergers lead to early-type galaxies, the combined effect of gas-rich and dry mergers is able to explain most of the evolution in the number density of massive early-type galaxies since $z \sim 1.5$, with our measured gas-rich merger rate accounting for about two-thirds of this evolution.

Conclusions. Merging of star-forming galaxies is frequent at around the peak in star formation activity. Our results show that gas-rich mergers make an important contribution to the growth of massive galaxies since $z \sim 1.5$, particularly on the build-up of the red sequence.

Key words. galaxies:evolution — galaxies:formation — galaxies:interactions

1. Introduction

Understanding the mechanisms involved in the mass assembly of galaxies and their relative role over cosmic time is an important open topic in modern astrophysics. In particular, the evolution of the red sequence, which includes passive galaxies dominated by old stellar populations and an early-type (E/S0) morphology,

imposes fundamental constraints on the formation and evolution models.

The stellar mass density in the red sequence has increased by a factor of ~ 10 in the 2.5 Gyr between $z = 2$ and $z = 1$, but only by a factor of ~ 2 in the last 7 – 8 Gyr of cosmic history (e.g., Arnouts et al. 2007; Vergani et al. 2008; Ilbert et al. 2010). Major mergers, the merger of two galaxies with similar stellar masses, is an efficient mechanism for creating new passive, early-type galaxies (e.g., Naab et al. 2006; Rothberg & Joseph 2006a,b; Hopkins et al. 2008; Rothberg & Fischer 2010; Bournaud et al. 2011). Thus, the knowledge of the merger rate at $z > 1$ is important input when estimating the relative contribution of merging and cold-gas accretion (e.g., Dekel & Birnboim 2006) in the early assem-

* This work is based mainly on observations collected at the European Southern Observatory (ESO) Very Large Telescope (VLT), Paranal, Chile, as part of the Programs 179.A-0823, 177.A-0837, 78.A-0177, 75.A-0318, and 70.A-9007. This work also benefits from data products produced at TERAPIX and the Canadian Astronomy Data Centre as part of the Canada-France-Hawaii Telescope Legacy Survey, a collaborative project of NRC and CNRS.

bly of galaxies and, in particular, the role of merging in the build-up of the red sequence.

The evolution of the merger rate since $z \sim 1$ is now well constrained by direct observations. The early measurements using photometric pairs (Patton et al. 1997; Le Fèvre et al. 2000) or post-merger morphological signatures (Conselice 2003; Jogee et al. 2009) have been superseded by spectroscopic measurements confirming physical pairs from the redshift measurement of both components of a major merger with a luminosity/mass ratio $\mu \geq 1/4$ (e.g., Lin et al. 2008; de Ravel et al. 2009, 2011), as well as for minor mergers down to $\mu = 1/10$ (López-Sanjuan et al. 2011). With a parametrization of the merger rate's evolution following $\propto (1+z)^n$, it is observed that the major merger rate's evolution depends on the luminosity and on the mass of the galaxy sample (e.g., de Ravel et al. 2009), where massive galaxies with $M_* > 10^{11} M_\odot$ have a higher merger rate, but with little redshift evolution ($n \sim 0 - 2$), while lower mass galaxies with $M_* = 10^9 - 10^{11} M_\odot$ have a lower merging rate but with stronger redshift evolution ($n \sim 3 - 4$). This mass dependency seems to explain some of the apparent discrepancy of merger rate measurements made from observations targeting different mass samples.

Beyond $z \sim 1$, direct measurements of the merger rate are still limited. Previous attempts to measure the major merger rate at $z > 1$ have focused on the identification of merger remnants from morphological signatures (Conselice et al. 2008, 2011b; Bluck et al. 2012), on the study of projected close pairs (Ryan et al. 2008; Bluck et al. 2009; Williams et al. 2011; Man et al. 2012; Mármol-Queraltó et al. 2012; Law et al. 2012), or on indirect estimations (Cameron & Pettitt 2012; Puech et al. 2012). These studies find a high merger rate to $z \sim 2 - 3$ but with a large scatter between different measurements. However, these results are up to now solely based on photometric measurements which are increasingly hard to correct for contamination along the line of sight as redshift increases. Another complication stems from the morphological evolution of galaxies, with show more irregular morphologies at high redshifts, and a wavelength dependency with more multi-component objects present when observed in the rest-frame UV (Law et al. 2007), with some of these components possibly related to strong star-forming regions rather than to different dynamical components.

To improve on this situation, it is necessary to obtain spectroscopic confirmation of the physical nature of the photometric pairs at $z \geq 1$. In the last years, NIR Integral Field Spectrographs (IFSs), like SINFONI on the VLT or OSIRIS on the Keck, have opened the possibility for a systematic study of the dynamical field around high redshift galaxies in the optical rest-frame. Some examples are the MASSIV¹ (Mass Assembly Survey with SINFONI in VVDS, Contini et al. 2012) survey at $0.9 < z < 1.8$, the SINS² (Spectroscopic Imaging survey in the Near-infrared with SINFONI, Förster Schreiber et al. 2009) survey at $z \sim 2$, or the Keck-OSIRIS (Law et al. 2009), the AMAZE (Assessing the Mass-Abundance redshift -Z- Evolution, Maiolino et al. 2008) and the LSD (Lyman-break galaxies Stellar populations and Dynamics, Mannucci et al. 2009) surveys at $z \sim 3$.

The MASSIV survey has been designed to target the peak of the star-formation rate at $0.9 < z < 1.8$, filling the gap between higher redshift ($z \sim 2$) IFS surveys with those at $z < 1$, e.g., IMAGES (Intermediate Mass Galaxy Evolution Sequence, Yang et al. 2008). The MASSIV survey has targeted 84 star-forming galaxies at $0.9 < z < 1.8$ with SINFONI, drawn

from the VVDS³ (VIMOS VLT Deep Survey, Le Fèvre et al. 2005) survey. MASSIV has been used as a unique opportunity to study in detail the dynamical state of $0.9 < z < 1.8$ galaxies (Epinat et al. 2012), their metallicity gradients (Queyrel et al. 2012), or the evolution of the fundamental mass-size-velocity relations since $z \sim 1.2$ (Vergani et al. 2012).

In this paper, using the MASSIV survey, we present for the first time a measurement of the gas-rich major merger rate of star-forming galaxies from kinematical close pairs at $0.9 < z < 1.8$. Thanks to the large field-of-view of IFS we have access to the complete surrounding volume of the galaxies when searching for close kinematical companions. In addition, the well-defined selection of sources from the VVDS and the well controlled selection function of MASSIV observations ensures the study of a representative population of star-forming galaxies at these redshifts (see Contini et al. 2012, for details). This all together enables the measurement of average volume quantities like the merger fraction and rate.

The paper is organised as follows: in Sect. 2 we summarise the MASSIV data set used to identify merging pairs, and in Sect. 3 we develop the methodology to measure the merger fraction from IFS data. We report the gas-rich major merger fraction in MASSIV in Sect. 4, and derive the gas-rich major merger rate in Sect. 5. We discuss the implication of our results in Sect. 6. Finally, we present our conclusions in Sect. 7. We use $H_0 = 100h$ km s⁻¹ Mpc⁻¹, $h = 0.7$, $\Omega_m = 0.3$, and $\Omega_\Lambda = 0.7$ throughout this paper. All magnitudes refer to the AB system. The stellar masses assume a Salpeter (1955) initial mass function (IMF).

2. The MASSIV data set

The galaxy sample studied in this paper is the final release of the MASSIV project (ESO Large Programme 179.A-0823; PI.: T. Contini). A full description of the sample can be found in Contini et al. (2012). We briefly summarise some properties of this sample of 84 galaxies below. The galaxies were selected from the VVDS in the RA = 2h area of the deep ($I_{AB} \leq 24$; Le Fèvre et al. 2005) and ultradeep ($I_{AB} \leq 24.75$; Le Fèvre et al., in prep.) surveys, and in the RA = 14h and RA = 22h areas of the wide survey ($I_{AB} \leq 22.5$; Garilli et al. 2008).

The MASSIV sources are a subsample of the VVDS star-forming population at $z > 0.9$. The star-forming selection was performed on the measured intensity of [O II] $\lambda 3727$ emission line in the VIMOS spectrum (see Lamareille et al. 2009; Vergani et al. 2008) or, for the cases where the [O II] $\lambda 3727$ emission line was out of the VIMOS spectral range (i.e., for $z \geq 1.5$ galaxies), on the UV flux based on their observed photometric *UBVR*IK spectral energy distribution and/or UV rest-frame spectrum. The star formation criteria ensure that rest-frame optical emission lines H α and [N II] $\lambda 6584$, or in a few cases [O II] $\lambda 5007$, will be bright enough to be observed with SINFONI in the NIR *J* (sources at $z < 1.1$) and *H* (sources at $z > 1.1$) bands. In addition, to ensure that the H α emission line will be detected with SINFONI with a sufficient signal-to-noise (*S/N*) ratio in a reasonable exposure time, an additional selection in the [O II] $\lambda 3727$ equivalent width of $0.9 < z < 1.5$ sources is imposed. The final selection of parent VVDS star-forming galaxies in the [O II] $\lambda 3727$ flux vs equivalent width plane at $0.9 < z < 1.5$ is shown in Fig. 1.

The 84 MASSIV sources were randomly selected from the star-forming population in the VVDS and also fulfil two important observational constraints, (i) the observed wavelength of

¹ <http://www.ast.obs-mip.fr/users/contini/MASSIV/>

² http://www.mpe.mpg.de/forster/SINS/sins_nmf.html

³ <http://cesam.oamp.fr/vvdsproject/>

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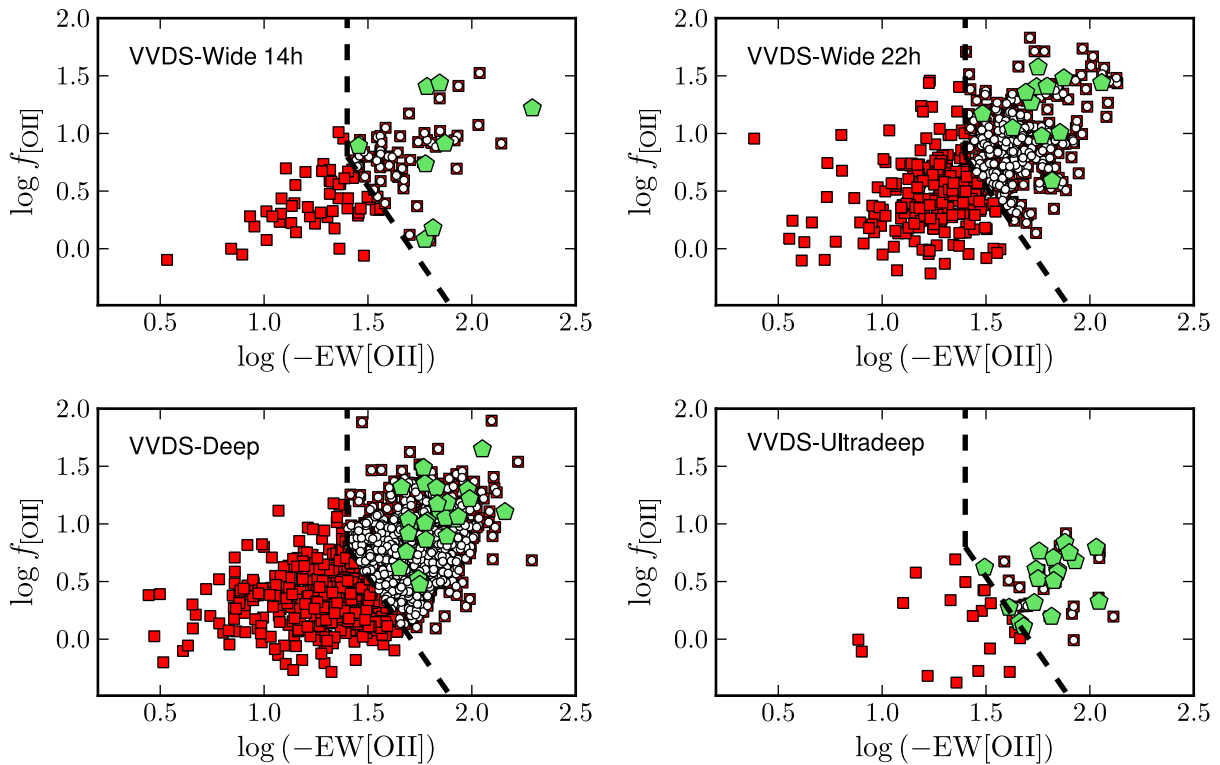


Fig. 1. MASSIV selection in the $[\text{O II}]\lambda 3727$ flux [10^{-17} erg s^{-1} cm^{-2}] vs equivalent width [\AA] plane (see Contini et al. 2012, for details). Red squares are the VVDS sources at $0.94 < z < 1.5$ with individual $[\text{O II}]\lambda 3727$ line measurement in the VVDS-Wide 14h (top-left), VVDS-Wide 22h (top-right), VVDS-Deep (bottom-left), and VVDS-Ultradeep surveys (bottom-right). The dashed lines mark the selection of MASSIV star-forming galaxies. White dots are those VVDS sources that fulfil the MASSIV selection. Green pentagons are the MASSIV galaxies observed with SINFONI/VLT. [A colour version of this plot is available at the electronic edition].

$\text{H}\alpha$ line falls 9\AA away from strong OH night-sky lines, in order to avoid heavy contamination of the galaxy spectrum by sky-subtraction residuals. And (ii) a bright star ($R < 18$ mag) is close enough to the target to observe it at higher spatial resolution with the adaptive optics (AO) system of SINFONI.

The most stringent selection criterion used to build the MASSIV sample is certainly the requirement of a minimum $[\text{O II}]\lambda 3727$ equivalent width. This sensitivity limit is likely to translate into an overall bias towards younger and more actively star-forming systems. To check for this possible effect, Contini et al. (2012) compare the properties of the MASSIV sample with those of the global VVDS sample, which is purely apparent magnitude-selected without any colour selection and thus representative of the overall star-forming population of galaxies at high redshifts. They conclude that the final MASSIV sample provides a good representation of “normal” star-forming galaxies at $0.9 < z < 1.8$ in the stellar mass regime $M_{\star} = 10^9 - 10^{11} M_{\odot}$, with a median star formation rate $SFR \sim 30 M_{\odot} \text{ yr}^{-1}$ and a detection limit of $\sim 5 M_{\odot} \text{ yr}^{-1}$.

The observations have been performed between April 2007 and January 2011. Most (85%) of the galaxies in the sample have been observed in a seeing-limited mode (with a spatial sampling of 0.125 arcsec/pixel). However, eleven galaxies have been acquired with AO assisted with a laser guide star (AO/LGS, seven with 0.05 and four with 0.125 arcsec/pixel spatial sampling). The data reduction was performed with the ESO SINFONI pipeline

(version 2.0.0), using the standard master calibration files provided by ESO. The absolute astrometry for the SINFONI data cubes was derived from nearby bright stars also used for point spread function (PSF) measurements. Custom IDL and Python scripts have been used to flux calibrate, align, and combine all the individual exposures. For each galaxy a non sky-subtracted cube was also created, mainly to estimate the effective spectral resolution. For more details on data reduction, we refer to Epinat et al. (2012).

We use the $\text{H}\alpha$ emission line (or $[\text{O III}]\lambda 5007$ in a few cases) in the SINFONI data cubes to derive the kinematical maps (flux, velocity field and velocity dispersion map) of the MASSIV galaxies. To estimate their dynamical properties, we assume that the ionised gas rotates in a thin disc with two regimes for the rotation velocity, a solid body shape in the innermost regions and a plateau in the outskirts. Using a χ^2 minimization we produce seeing-corrected velocity and dispersion maps of galaxies with geometrical inputs weighted for the S/N ratio of each pixel. We estimate the geometrical parameters used in the fitting model on the i -band best-seeing CFHTLS Megacam images for all galaxies (Goranova et al. 2009), except for VVDS-Wide 14h galaxies that were covered with the CFHT-12K/CFHT camera (Le Fèvre et al. 2004). We use the GALFIT software (Peng et al. 2002) that convolves a PSF with a model galaxy image based on the initial parameter estimates fitting a Sérsic (1968) profile. Residual maps from the fitting were used to optimise the results.

At the end of the fitting procedure, GALFIT converges into a final set of parameters such as the centre, the position angle, and the axial ratio. The i -band images were also used to correct for SINFONI astrometry, using the relative position of the PSF star. The morphology and kinematics maps of the 50 first-epoch galaxies are presented in Epinat et al. (2012) together with a more extensive discussion on the model fitting procedure. The 34 second-epoch galaxies, already included in the present work, will be presented in a future paper.

We obtain the stellar mass of MASSIV galaxies from a spectral energy distribution (SED) fit to the photometric and spectroscopic data with Bruzual & Charlot (2003) stellar population synthesis models using the GOSSIP2 Spectral Energy Distribution tool (Franzetti et al. 2008). We assume a Salpeter (1955) IMF, and a set of delayed exponential star formation histories with galaxy ages in the range from 0.1 to 15 Gyr. As input for the SED fitting, in addition to the VVDS spectra, we use the multi-band photometric observations available in the VVDS fields (see Contini et al. 2012, for further details). Following Walcher et al. (2008) we adopt the probability distribution function to obtain the stellar mass.

In all, the final 84 MASSIV galaxies are representative of the normal star-forming ($SFR \gtrsim 5M_{\odot} \text{ yr}^{-1}$) population of $M_{\star} = 10^9 - 10^{11} M_{\odot}$ galaxies at $0.9 < z < 1.8$.

3. Measuring the merger fraction from IFS data

We define as a close pair two galaxies with a projected separation in the sky plane $r_p^{\min} \leq r_p \leq r_p^{\max}$ and a rest-frame relative velocity along the line of sight $\Delta v \leq \Delta v^{\max}$. We used $\Delta v^{\max} = 500 \text{ km s}^{-1}$ and $r_p^{\max} = 20 - 30 h^{-1} \text{ kpc}$ (see Sect. 3.3, for details), while setting $r_p^{\min} = 0$. We therefore searched for close companions in the kinematical maps of the MASSIV sources, analysing and classifying the sample using the velocity field and the velocity dispersion map (see Epinat et al. 2012, for details about the classification). Note that spectroscopic redshifts from the VVDS sources outside MASSIV are not used in this close pair search and only those sources detected in the SINFONI data cubes are taken into account. We find 20 close pair candidates in the MASSIV data cubes, and we study these systems in detail to select major (luminosity difference between both components $\mu = L_2/L_1 \geq 1/4$) close pairs (Sect. 4).

If there was N_p major close pairs in our sample, the major merger fraction is

$$f_{\text{MM}} = \frac{N_p}{N}, \quad (1)$$

where N is the number of principal galaxies targeted for the survey. We named principal galaxy the source in the pair closest to the kinematical centre of the targeted system, even if it is not the brightest/more massive galaxy in the pair. In addition, we assume that all our close pairs are gas-rich: as shown by Vergani et al. (2012), the gas fraction of the MASSIV sources is above 10%, with a median value of $\sim 30\%$. The simple definition in Eq. (1) is only valid for volume-limited samples. While our sample is not only luminosity-limited but spectroscopically defined, we must take into account the different selection effects, both in the VVDS parent samples and in MASSIV, in our computation of the merger fraction.

3.1. Accounting for selection effects in the VVDS

Since a fraction of the total number of potential targets in the VVDS fields have been spectroscopically observed and since

the redshifts are not measured with 100% accuracy, we must correct for the VVDS Target Sampling Rate (TSR) and the Redshift Success Rate (SSR), computed as a function of redshift and source magnitude. The SSR has been assumed independent of galaxy type, as demonstrated up to $z \sim 1$ in Zucca et al. (2006). Every VVDS source has a redshift confidence flag (see Le Fèvre et al. 2005, for details), that can be flag = 4 (redshift 99% secure), flag = 3 (97% secure), flag = 2 (87% secure), flag = 9 (redshift from a single emission line, 90% secure), flag = 1 (50% secure), or flag = 0 (no redshift information). As several VVDS-Deep galaxies with flag = 2 have been re-observed in the VVDS-Ultradeep survey, providing a robust measurement of their redshift, this offers the opportunity to estimate the reliability of VVDS-Deep flag = 2 sources. We thus define a weight w_{29} to take this into account. We also define the weight w_{29} for flag = 9 sources by comparison with the latest photometric redshifts in the VVDS-Deep field (see Cucciati et al. 2010, for details about the latest photometric data set in the 2h field). By definition, the w_{29} weight is equal to 1 for flag = 3 and 4 sources in VVDS-Deep, and for all sources in VVDS-Wide and VVDS-Ultradeep. We derived the spectroscopic completeness weight for each galaxy i in the VVDS catalogue as

$$w_{\text{VVDS}}^i(z, I_{\text{AB}}, \text{flag}) = \frac{w_{29}^i(z, I_{\text{AB}}, \text{flag})}{TSR^i(z, I_{\text{AB}}) \times SSR^i(z, I_{\text{AB}})}. \quad (2)$$

The TSR , SSR and w_{29} on VVDS-Deep and VVDS-Ultradeep were measured in previous works (Ilbert et al. 2006; Cucciati et al. 2012). We assume $TSR = 0.22$ in VVDS-Wide fields (Garilli et al. 2008), and we detail the computation of the SSR in VVDS-Wide fields in Appendix A.

3.2. Accounting for selection effects in the MASSIV survey

The MASSIV sources were randomly drawn from the star-forming population in the VVDS (Sect. 2). We correct for three basic selection effects in MASSIV.

- The selection weight, w_{sel} . We define this weight as the fraction of star-forming galaxies that fulfil the MASSIV selection in the $[\text{O II}]\lambda 3727$ flux [$10^{-17} \text{ erg s}^{-1} \text{ cm}^{-2}$] vs equivalent width [\AA] plane at $z < 1.5$ (Fig. 1). The weight w_{sel} tells us how representative the MASSIV selection of the global star-forming population in each of the VVDS surveys (Wide, Deep and Ultradeep) is, and gives more importance to the more representative samples. The VVDS-Deep and Ultradeep have $w_{\text{sel}} \sim 0.67$ at $z < 1.5$, while the VVDS-Wide fields have $w_{\text{sel}} \sim 0.53$ in the same redshift range. We assume $w_{\text{sel}} = 1$ at $z \geq 1.5$, where the selection is based on colour/spectral properties and all the star-forming galaxies in the VVDS are thus pre-selected;
- the MASSIV IFS Rate (MIR) is defined as the fraction of galaxies that fulfil the MASSIV selection and which were finally observed with SINFONI (Fig. 1). The MIR ranges from 0.43 for VVDS-Ultradeep to 0.05 for VVDS-Deep;
- the MASSIV Success Rate (MSR) is the fraction of observed sources with a reliable kinematical classification. This fraction is always high, $MSR \gtrsim 0.8$.

Finally, the MASSIV weight is

$$w_{\text{MASSIV}}^j(x, z) = \frac{w_{\text{sel}}(x, z)}{MIR^j(x, z) \times MSR^j(x, z)}, \quad (3)$$

where the index j spans for the MASSIV sources and x refers to the VVDS survey (Wide at 14h, Wide at 22h, Deep or Ultradeep) to which the source belongs.

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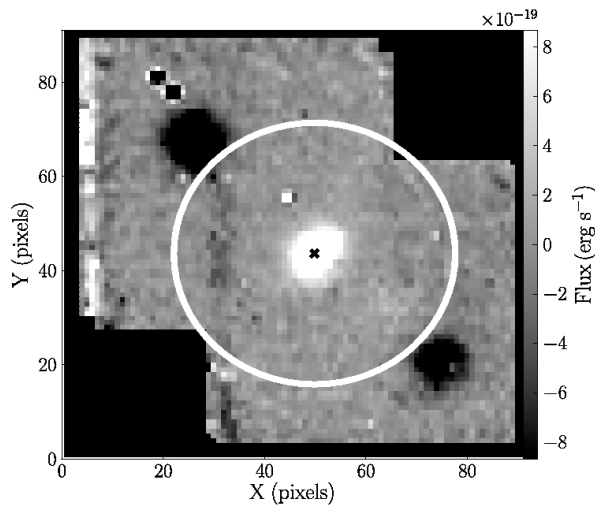


Fig. 2. Typical field-of-view of a MASSIV source. We show the flux of the source 220397579 in the five channels of its reduced, flux calibrated, data cube centred on the position of the $H\alpha$ emission line (scale in the right). The pixel spatial scale is $0.125''$, that at the redshift of the source ($z = 1.0379$) corresponds to ~ 1 kpc. The source is at the centre of the image (black cross), while the white circle delimits the $20h^{-1}$ kpc area around the source. In this particular case, 99% of this area is covered with MASSIV data and the two negatives of the source (black regions) are also excluded.

3.3. Area correction

Thanks to the large field-of-view of SINFONI, we have access to the complete surrounding volume of the principal galaxy when searching for close companions. This is a major advantage with respect to long-slit spectroscopic surveys, in which the observed number of close pairs is proportional to TSR^2 , diminishing the statistics. However, the final reduced MASSIV data cubes cover a finite area in the sky plane. To deal with this, we define the area weight as

$$w_{\text{area}}^j(z, r_p^{\text{max}}) = \frac{A_r}{A_{\text{MASSIV}}}, \quad (4)$$

where A_r is the area subtended in the sky plane by a circle of radius r_p^{max} at the redshift of the source j , and A_{MASSIV} is the area of the same circle covered by the reduced SINFONI mosaic (Fig. 2). We state that the optimum search radius is $r_p^{\text{max}} = 20h^{-1}$ kpc. This choice of r_p^{max} minimises the area correction, with $w_{\text{area}} \sim 1$ in all cases. In addition, the two negatives of the main source produced by the offset observing procedure and that appear in the extremes of the reduced SINFONI mosaic are then excluded from the search area.

Finally, the corrected gas-rich major merger fraction is

$$f_{\text{MM}} = \frac{\sum_k^{N_p} w_{\text{VVDS}}^k w_{\text{MASSIV}}^k w_{\text{area}}^k}{\sum_j^N w_{\text{VVDS}}^j w_{\text{MASSIV}}^j}, \quad (5)$$

where the index j and k spans respectively for all MASSIV galaxies and for the MASSIV galaxies with a major close companion. The error budget in the major merger fraction is dominated by the low statistics. We use the Bayesian approach from Cameron (2011) to measure the statistical error in the raw major

merger fraction, i.e., in N_p/N , then scale it with the weighting scheme above.

4. Mergers classification and the gas-rich major merger fraction in MASSIV

In this section we measure, for the first time, the major merger fraction at $0.9 < z < 1.8$ from spectroscopically-confirmed close pairs. The MASSIV observational strategy defines three natural redshift bins (Sect. 2). The low redshift MASSIV sources were observed in the J band, while the higher redshift ones in the H band. This translates to a gap in the redshift distribution at $z \sim 1.1$. In addition, the selection function of MASSIV targets changes at $z = 1.5$, providing another redshift boundary. We take advantage of these natural separations in the data to estimate the gas-rich major merger fraction in three redshift ranges, $z_{r,1} = [0.94, 1.06]$, $z_{r,2} = [1.2, 1.5]$, and $z_{r,3} = [1.5, 1.8]$. We restrict our study to those galaxies with $I_{\text{AB}} \leq 23.9$ to ensure completeness in the detection of close pairs (see Appendix C. for details).

We follow the steps described below to split the close pairs candidates in the MASSIV data cubes into major and no major mergers:

1. As described in Epinat et al. (2012), we had performed a classification of the MASSIV sources based on the shape of the velocity field (regular or irregular) and the close environment (isolated or not-isolated). From this classification, we had pre-selected as close pair candidates the non-isolated sources and those identified as mergers from the velocity field (see Epinat et al. 2012, for more details). We identified 20 close pair candidates (Table 1).
2. To study in detail these close pair candidates we used the deepest i -band images in the VVDS fields: CFTH12K (14h-field, exposure time of $t_{\text{exp}} = 3.6$ ks, Le Fèvre et al. 2004), CFHTLS-Wide (22h field, $t_{\text{exp}} = 4 - 10$ ks, Goranova et al. 2009), and CFHTLS-Deep (02h field, $t_{\text{exp}} \sim 300$ ks, Goranova et al. 2009). We run SExtractor on the systems with well separated sources, and GALFIT (with two Sérsic components) on the blended ones, to estimate the luminosity difference in the i band between both sources, $\Delta m_i = m_{i,2} - m_{i,1}$. We took $\Delta m_i \leq 1.5$ (factor four or less in luminosity) to identify major mergers. The observed i band corresponds to $\sim 300 - 350$ nm rest-frame in the redshift range of our sample. We stress that for the blended sources, we run GALFIT v3.0 (Peng et al. 2010) without imposing any constraint to the parameters of the fit and we only used the information from the kinematical maps to set the initial positions of the sources. We present the residual maps of these blended sources in Appendix B.
3. We confronted the images and the two component fits from GALFIT with the velocity field and the velocity dispersion map of the sources. We compared the distribution of $H\alpha$ emission and the geometry of the velocity field to the rest-frame UV continuum (or rest-frame visible when NIR images are available). The presence of two components with position and geometry concordant in the velocity field and in the continuum images, is a strong indication of the reality of the pair.
4. We run SExtractor in the residual image from GALFIT with the principal source subtracted to obtain a second estimation of $m_{i,2}$, while with the companion source subtracted to estimate $m_{i,1}$. Then we compared the Δm_i derived from the GALFIT modeling with that from these SExtractor estimations. We found good agreement between both measure-

Table 1. Close pair candidates at $0.9 < z < 1.8$ in the MASSIV sample.

ID	z	r_p (h^{-1} kpc)	Δv (km s^{-1})	Δm_i	Δm_{K_s}	Classification
020294045	1.0028	2.9	180	0.4	0.3	Major merger
020386743	1.0487	3.7	80	1.8	...	No major merger
020461235	1.0349	2.8	15	1.7	...	No major merger
140096645	0.9655	> 1.5	...	No major merger
220397579	1.0379	14.4	340	0.4	-1.4	Major merger
220544394	1.0101	7.1	50	1.3	...	Major merger
020167131	1.2246	15.2	130	0.2	0.1	Major merger
020218856	1.3103	> 1.5	...	No major merger
020240675	1.3270	> 1.5	...	No major merger
020283083	1.2818	3.8	5	0.7	...	Major merger
020283830	1.3949	8.5	500	1.9	...	No major merger
020465775	1.3583	3.6	40	0.7	...	Major merger
220376206	1.2445	13.4	400	2.4	...	No major merger
220544103	1.3973	5.0	75	-1.1	...	Major merger
910154631	1.3347	4.2	130	0.8	...	Major merger
910296626	1.3558	12.1	165	-0.1	-0.2	Major merger
910337228	1.3955	9.5	220	1.4	...	Major merger
020116027	1.5302	26.8	100	0.7	0.5	Major merger
910186191	1.5399	12.7	450	-0.2	-2.4	Major merger
910274060	1.5694	3.4	10	0.2	...	Major merger

ments (difference of ~ 0.2 mag or less). The major merger classification did not change from the initial estimate.

- Finally, we also explored Δm_{K_s} for well separated galaxies and for one blended source (020294045). We used the K_s -band images from UKIDSS-DXS survey (22h field, Dye et al. 2006) and WIRDS (2h field, Bielby et al. 2012). The observed K_s band corresponds to $\sim 0.8-1 \mu\text{m}$ rest-frame in the redshift range of our sample, and is a better tracer of the stellar mass content of the galaxies. After this second check there is not change in the major merger classification, except for one close pair (source 910186191), supporting the previous i -band results.

We converged to the steps above after exploring different possibilities. The stellar mass ratio between the two galaxies in a close pair is the best parameter to classify such system as a major merger. However, the estimation of the stellar mass in the blended sources and in some well separated pairs is not feasible due to spatial resolution and photometric depth limitations. We concluded that the luminosity difference in the i band is an homogeneous criterion applicable to the whole MASSIV sample and have additional benefits (i) the observed i band corresponds to the UV continuum of the source, which is related with the star formation of the galaxy and thus with its $H\alpha$ emission. Hence, both pieces of information should provide a consistent picture about the system under study. (ii) The MASSIV PSF is $0.5 - 0.8''$, similar to the typical seeing in the i band, $\sim 0.74''$, minimising spatial resolution differences. (iii) The 3σ detection magnitude of the CFTLHS i -band images in the 2h field is ~ 26 (AB). Thanks to this depth we are able to detect the outskirts of the fainter MASSIV galaxies, making feasible the decomposition of the blended sources. We tried the two components fit in redder bands, but the lower S/N lead in general to poor constraints. And (iv) the VVDS parent samples are i -band magnitude selected, so the completeness of the MASSIV close pair sample is well defined in the i band (see Appendix C, for further

details). In conclusion, even if Δm_i is not the optimal parameter to select major merger systems, it is the best practical one with the current data sets.

In the next sections we present the decomposition and the classification of the 20 close pair candidates in the MASSIV data cubes (Table 1).

4.1. Close pair candidates at $0.94 \leq z \leq 1.06$

The weighted mean redshift of the first redshift bin is $\bar{z}_{r,1} = 1.03$. We have identified 6 close pairs over 18 sources in this redshift range:

- 020294045 (Fig. 3). **Major merger.** The velocity map suggests two projected components. The companion is toward the north and presents a steep velocity gradient compared with the principal galaxy. The GALFIT model with two components in the i band reproduces both the position of the two kinematical components in the velocity map and the high velocity dispersion in the overlapping region between both components. The luminosity difference is $\Delta m_i = 0.4$. The separation between the sources is $2.9h^{-1}$ kpc and their relative velocity is $\Delta v \sim 180 \text{ km s}^{-1}$. Despite the close separation lead into significant overlap, this system also gives a satisfactory fit in the WIRDS K_s -band image (Fig.4). In this case $\Delta m_{K_s} = 0.3$, confirming that this system is a major merger.
- 020386743 (Fig. 5). *No major merger.* The velocity map suggests two projected components. The companion is toward the north and presents a different velocity than the northern part of the principal galaxy. The GALFIT model with two components suggests that we are only detecting the western part of the companion galaxy, with the eastern part being too faint. This is also consistent with the velocity dispersion map, which shows a regular pattern in the overlapping region. The separation between the sources is $3.7h^{-1}$ kpc and

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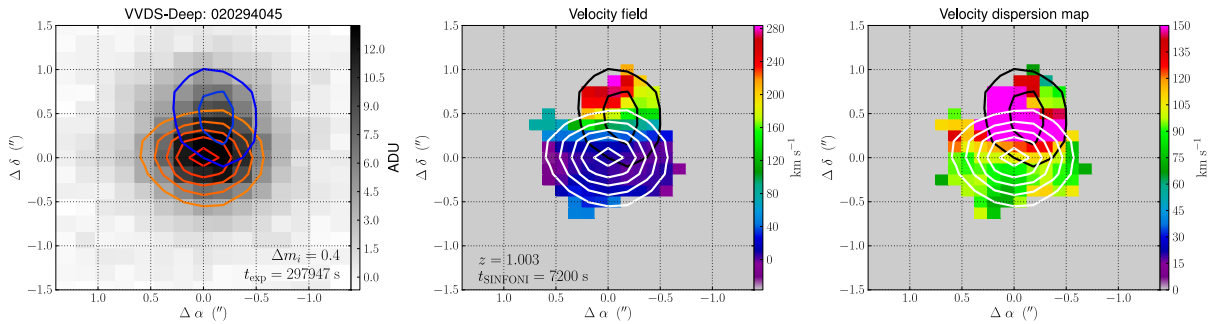


Fig. 3. The i -band image (*left panel*, scale in ADU in the right), the velocity field (*central panel*, scale in km s^{-1} in the right), and the velocity dispersion map (*right panel*, scale in km s^{-1} in the right) of the MASSIV source 020294045 (major merger). On each map, North is up and East is left. The level contours mark the isophotes of the two components obtained with GALFIT in the i -band image. The principal galaxy (red/white) is the one closer to the kinematical centre of the system and sets the origin in right ascension (α) and declination (δ), while the companion (blue/black) is the secondary component. The outer contour marks the 3.43 ADU ($7\sigma_{\text{sky}}$) isophote. The next contours mark brighter isophotes in 1.96 ADU ($4\sigma_{\text{sky}}$) steps. The luminosity difference between both components in the i band, Δm_i , as well as the CFHTLS exposure time, t_{exp} , are shown in the left panel. The redshift of the source and the on-source SINFONI exposure time, t_{SINFONI} , are shown in the central panel. [A colour version of this plot is available at the electronic edition].

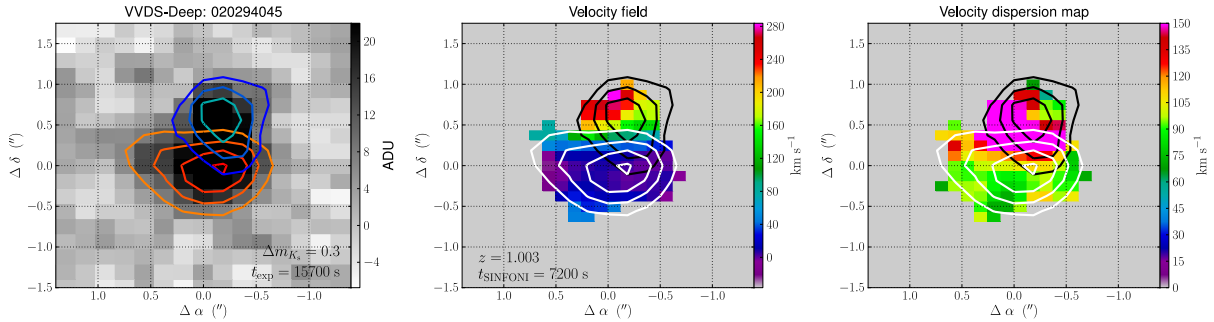


Fig. 4. The same as Fig. 3, but for the K_s -band image (*left panel*) of the MASSIV source 020294045 (major merger). The outer contour marks the 6.76 ADU ($2\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 5.07 ADU ($1.5\sigma_{\text{sky}}$) steps. The flux difference between both components in the K_s band, Δm_{K_s} , as well as the WIRDS exposure time, t_{exp} , are shown in the left panel. [A colour version of this plot is available at the electronic edition].

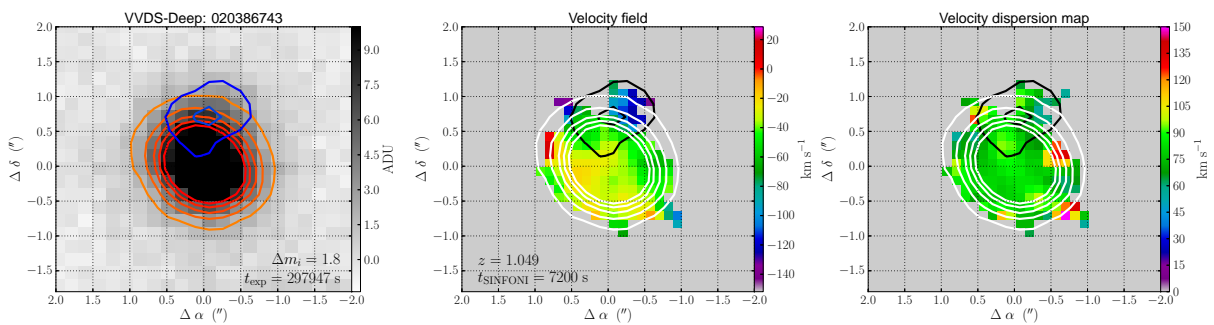


Fig. 5. The same as Fig. 3, but for the MASSIV source 020386743 (no major merger). The outer contour marks the 1.41 ADU ($2\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 1.76 ADU ($2.5\sigma_{\text{sky}}$) steps up to 10 ADU to avoid crowded figures. [A colour version of this plot is available at the electronic edition].

the luminosity difference is $\Delta m_i = 1.8$. Hence, we do not classify the system as a major merger.

- 020461235 (Fig. 6). *No major merger.* The velocity map suggests two projected components. The companion is to-

ward the south-west and presents a nearly constant velocity, in contrast with the velocity gradient of the principal galaxy. The GALFIT model with two components suggests that the companion is highly distorted, perhaps because the system is

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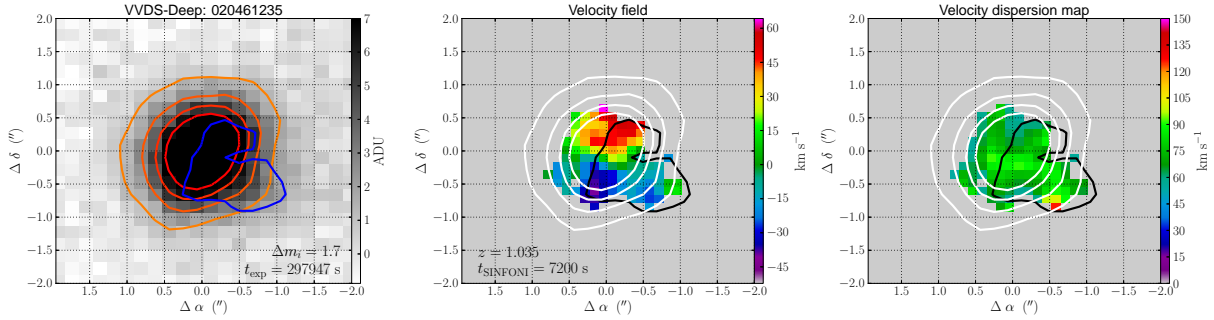


Fig. 6. The same as Fig. 3, but for the MASSIV source 020461235 (no major merger). The outer contour marks the 1.31 ADU ($3\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 1.75 ADU ($4\sigma_{\text{sky}}$) steps up to 7 ADU to avoid crowded figures. [A colour version of this plot is available at the electronic edition].

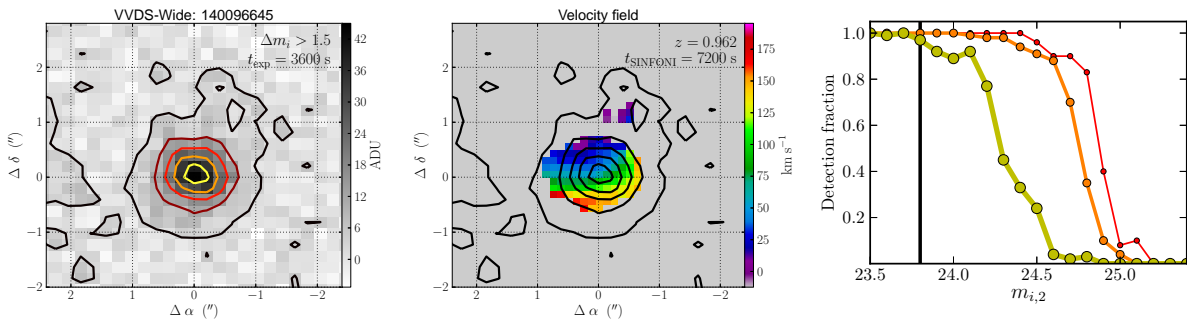


Fig. 7. Left and central panels are the same as in Fig. 3, but for the MASSIV source 140096645 (no major merger). The level contours mark the isophotes of the original i -band image. The $H\alpha$ companion source is not detected in the i band. The outer contour marks the 2.64 ADU ($1\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 7.94 ADU ($3\sigma_{\text{sky}}$) steps up to 45 ADU to avoid crowded figures. The right panel shows the detection fraction of fake galaxies injected in the i -band image as a function of the magnitude of the companion source $m_{i,2}$. The size of the dots marks the detection curves for extended, normal and compact galaxies (see text for details). The vertical line marks the limiting magnitude for major companion ($\Delta m_i = 1.5$). [A colour version of this plot is available at the electronic edition].

in an advanced merger stage (i.e., after first pericenter passage or pre-coalescence). The luminosity difference from the GALFIT model is $\Delta m_i = 1.7$, while from the residual maps is $\Delta m_i = 1.9$. This difference strongly suggests that this system is not a major merger. The separation between the components is $2.8h^{-1}$ kpc.

- 140096645 (Fig. 7). *No major merger.* A small companion is identified in the velocity map about 1.2 arcsec away from the principal galaxy. There is no continuum detection in the i -band image in the position of the $H\alpha$ companion, although there is a $\sim 1\sigma_{\text{sky}}$ excess emission. To estimate if the non-detected companion could be bright/massive enough to lead to a major merger, we injected fake companion sources in the i -band image and studied their detection fraction as a function of the fake companion luminosity $m_{i,2}$. The fake sources were modelled with a Sérsic function and convolved with a typical PSF of the 14h field using GALFIT. For each fake source we assumed a random inclination and position angle, a Sérsic index $n_s = 1$, i.e., an exponential disc (we checked that the detection curve is similar assuming either $n_s = 0.5$ or $n_s = 2$), and an effective radius r_e given by the $r_e - M_\star$ relation in MASSIV, $\log r_e = 0.36 + 0.37[\log(M_\star/M_\odot) - 10]$ (see also Vergani et al. 2012). To estimate the stellar mass of the fake companion, we took Δm_i as a proxy of the mass ra-

tio between the principal galaxy, for which the stellar mass is known, and the companion. Then, we applied Poissonian noise to the model and injected it in the expected position of the possible companion. We measured, for different luminosities $m_{i,2}$, which fraction of the 500 injected fake sources were detected. We repeated the previous steps for fake sources with $2r_e$ (extended sources) and $0.5r_e$ (compact sources), spanning all the possible sizes of the real sources. We show the result of this experiment in the right panel of Fig. 7.

We find that even for extended sources, we are $\sim 90\%$ complete at $m_{i,2} \sim 24$, while the limiting magnitude for a major companion is $m_{i,2}^{\text{MM}} = 23.8$ (vertical line in the right panel of Fig. 7). Because of the small probability of non detection of a major companion, we classify this system as no major merger.

- 220397579 (Fig. 8). **Major merger.** The velocity map shows two different components with a separation of $r_p = 14.4h^{-1}$ kpc and a relative velocity of $\Delta v \sim 340$ km s^{-1} . The companion is toward the north-west. The luminosity difference is $\Delta m_i = 0.4$, suggesting a major merger. The difference in the K_s band is $\Delta m_{K_s} = -1.4$. The negative sign implies that the companion is more luminous than the principal, which is brighter in the NUV rest-frame and in $H\alpha$. This is consistent

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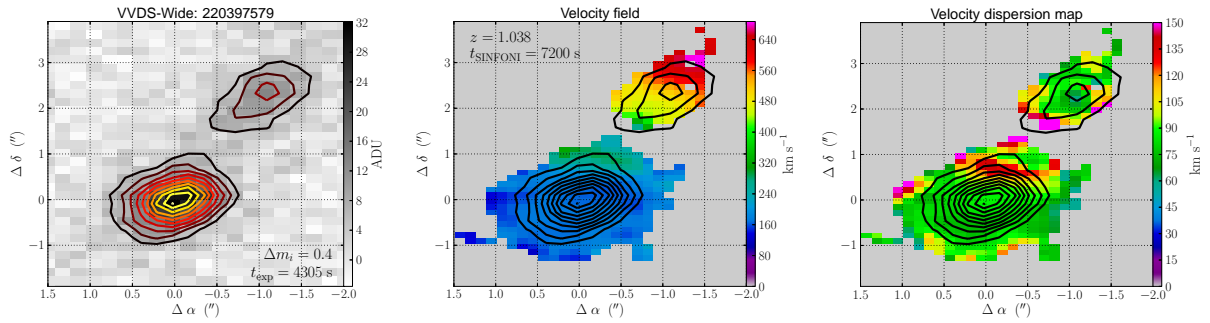


Fig. 8. The same as Fig. 3, but for the MASSIV source 220397579 (major merger). The level contours mark the isophotes of the original i -band image. The outer contour marks the 5.36 ADU ($3\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 2.68 ADU ($1.5\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

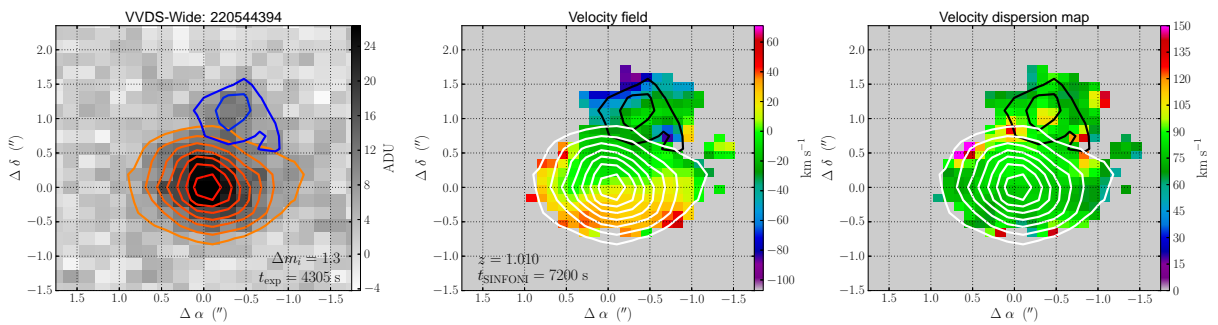


Fig. 9. The same as Fig. 3, but for the MASSIV source 220544394 (major merger). The outer contour marks the 4.31 ADU ($2\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 3.23 ADU ($1.5\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

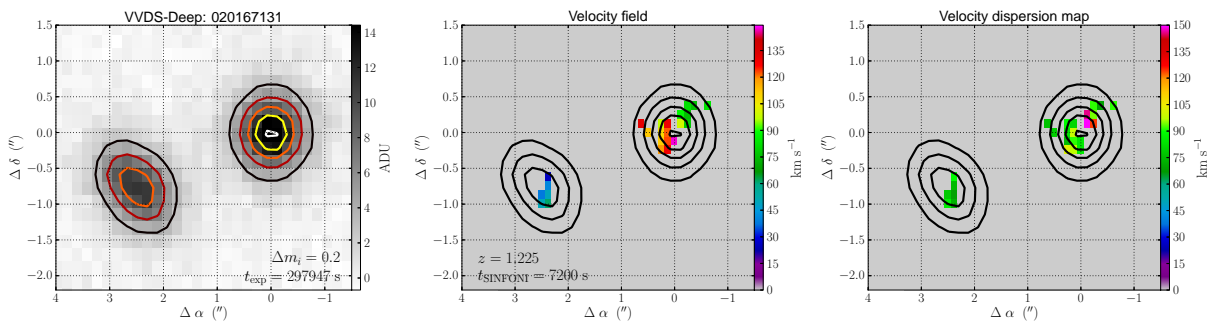


Fig. 10. The same as Fig. 8, but for the MASSIV source 020167131 (major merger). The outer contour marks the 3.43 ADU ($10\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 2.74 ADU ($8\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

with the measured integrated metallicity of these sources, that is higher for the companion (Queyrel et al. 2012). The suggested picture is that the system comprises a nearly face-on principal galaxy with intense star formation and low dust reddening, and a nearly edge-on companion galaxy with either a low level of star formation or strong dust reddening. We classify the system as a major merger.

- 220544394 (Fig. 9). **Major merger.** The velocity map shows two different components with a separation of $r_p = 7.1h^{-1}$ kpc and a relative velocity of $\Delta v \sim 50 \text{ km s}^{-1}$, that are also well recovered by the GALFIT model with two components.

The companion is toward the north, while the luminosity difference is $\Delta m_i = 1.3$. We classify the system as a major merger.

In summary for this redshift range, we classify 3 of the 6 close pair candidates as major mergers. This translates to a gas-rich major merger fraction of $f_{\text{MM}} = 0.208^{+0.152}_{-0.068}$ at $\bar{z}_{r,1} = 1.03$.

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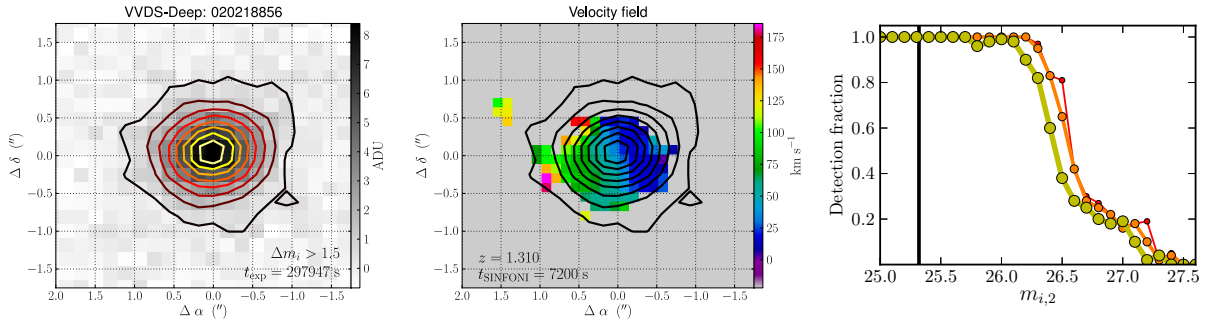


Fig. 11. The same as Fig. 7, but for the MASSIV source 020218856 (no major merger). The H α companion galaxy is not detected in the i band. The outer contour marks the 0.34 ADU ($1\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 1.01 ADU ($3\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

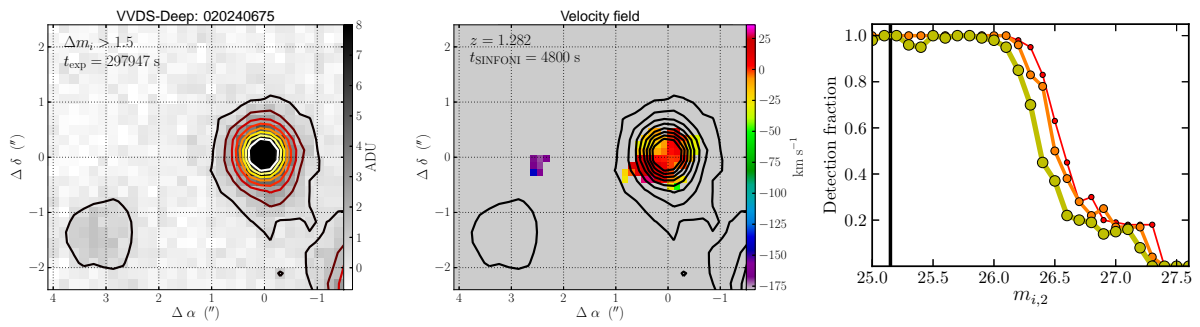


Fig. 12. The same as Fig. 7, but for the MASSIV source 020240675 (no major merger). The H α companion galaxy is not detected in the i band. The outer contour marks the 0.36 ADU ($1\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 1.08 ADU ($3\sigma_{\text{sky}}$) steps up to 8 ADU to avoid crowded figures. [A colour version of this plot is available at the electronic edition].

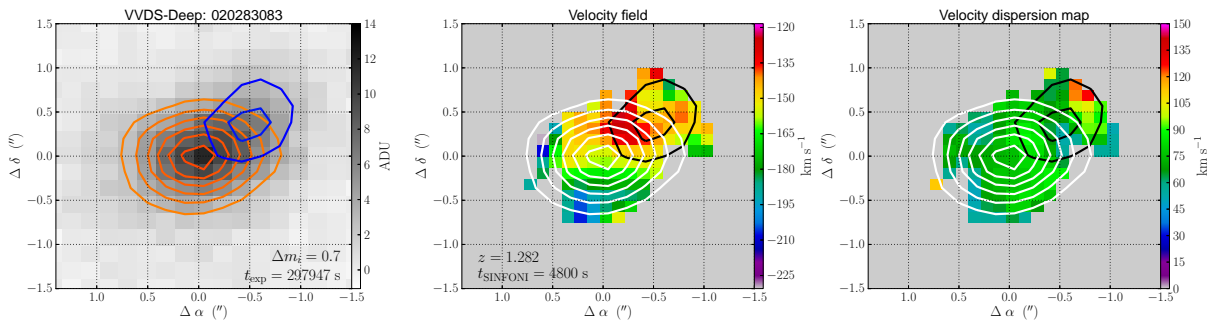


Fig. 13. The same as Fig. 3, but for the MASSIV source 020283083 (major merger). The outer contour marks the 2.12 ADU ($4\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 1.06 ADU ($2\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

4.2. Close pair candidates at $1.2 \leq z < 1.5$

The weighted mean redshift of the second redshift bin is $\bar{z}_{r,2} = 1.32$. This is a redshift range where there is no measurement of the major merger fraction from spectroscopic close pairs yet. In this framework, MASSIV provides an unique opportunity to measure the major merger fraction at this crucial epoch of galaxy evolution. We identify 11 close pair candidates over 30 galaxies in this redshift bin:

- 020167131 (Fig. 10). **Major merger.** The velocity map shows two different components with a separation of $r_p =$

$15.2h^{-1}$ kpc and a relative velocity of $\Delta v \sim 130 \text{ km s}^{-1}$. The line targeted in this case is [O III] $\lambda 5007$, which explains the low signal. The companion galaxy is toward the south-east, while the luminosity differences are $\Delta m_i = 0.2$ and $\Delta m_{K_s} = 0.1$. We classify the system as a major merger.

- 020218856 (Fig. 11). **No major merger.** A faint companion is detected in the velocity map about 1.6 arcsec from the principal galaxy. There is no continuum detection of this companion at the depth of the i -band image. We followed the same steps than for the source 140096645 to estimate the detection probability of companion galaxies, finding that we are

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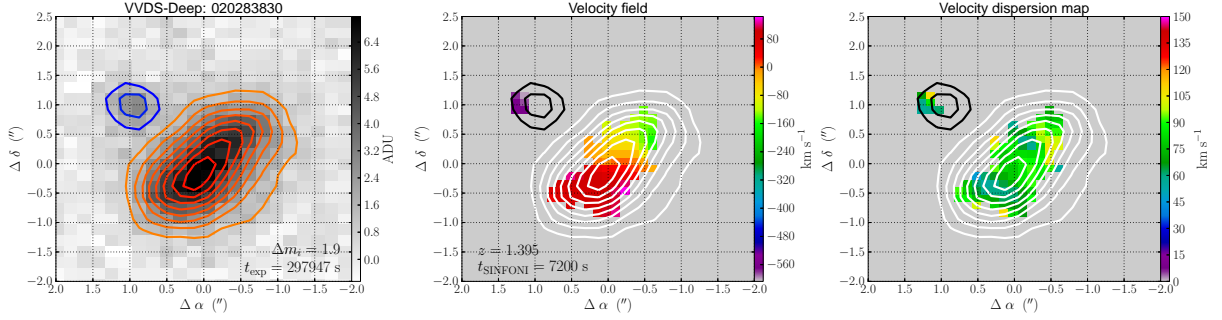


Fig. 14. The same as Fig. 3, but for the MASSIV source 020283830 (no major merger). The outer contour marks the 1.14 ADU ($3.5\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 0.75 ADU ($2.3\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

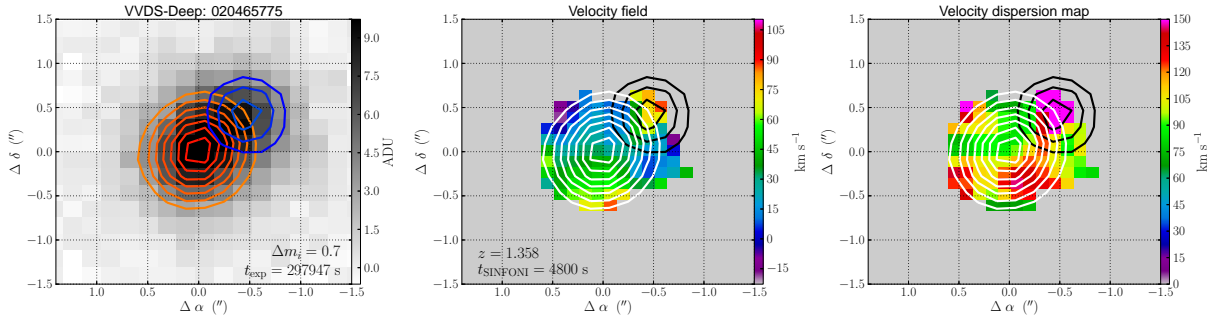


Fig. 15. The same as Fig. 3, but for the MASSIV source 020465775 (major merger). The outer contour marks the 2.92 ADU ($9\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 0.97 ADU ($3\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

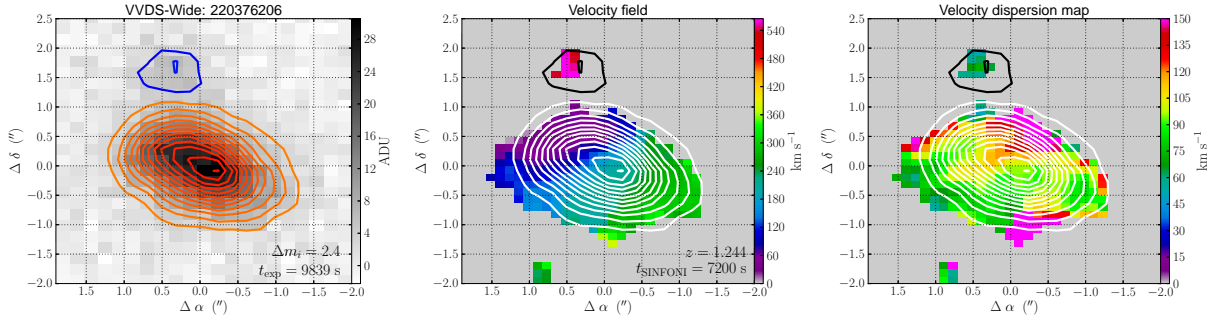


Fig. 16. The same as Fig. 3, but for the MASSIV source 220376206 (no major merger). The outer contour marks the 3.39 ADU ($3\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 2.26 ADU ($2\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

100% complete up to $m_{i,2} = 26$. We note that the curves for extended, normal and compact galaxies are similar, reflecting that the size of the PSF is larger than the assumed size of the fake sources. Because a major companion should be brighter than $m_{i,2}^{\text{MM}} = 25.32$ (vertical line in the left panel of Fig. 11), and the non detection in the continuum means that the companion should have $\Delta m_i \geq 2$, we do not classify the system as a major merger.

- 020240675 (Fig. 12). *No major merger.* A faint companion is detected in the velocity map about 2.4 arcsec from the principal galaxy. There is no continuum detection of this compan-

ion at the depth of the i -band image. We followed the same steps than for the source 020218856 to estimate the detection curves of companion galaxies. We find similar completeness curves. Because a major companion should be brighter than $m_{i,2}^{\text{MM}} = 25.15$ (vertical line in the left panel of Fig. 12), and the non detection in the continuum means that the companion should have $\Delta m_i \geq 2.5$, we do not classify the system as a major merger.

- 020283083 (Fig. 13). **Major merger.** The velocity map suggests two projected components separated by $3.8h^{-1}$ kpc and $\sim 5 \text{ km s}^{-1}$, with an extended region in the north-west. The

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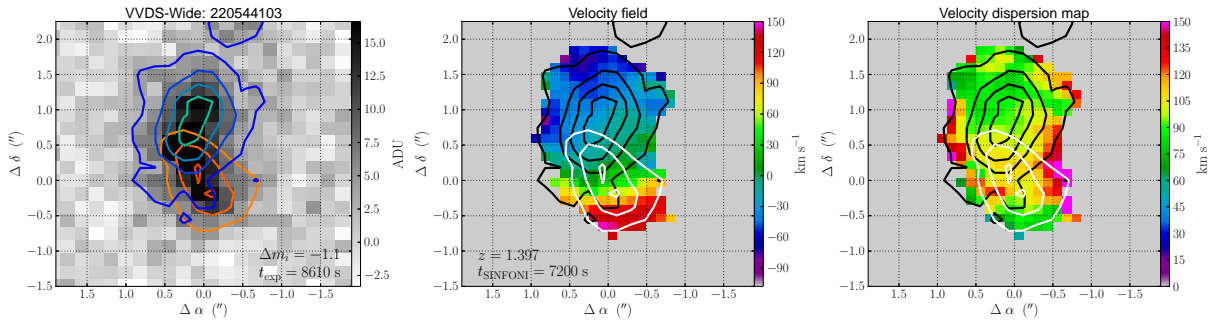


Fig. 17. The same as Fig. 3, but for the MASSIV source 220544103 (major merger). The outer contour marks the 3.32 ADU ($2\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 3.32 ADU ($2\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

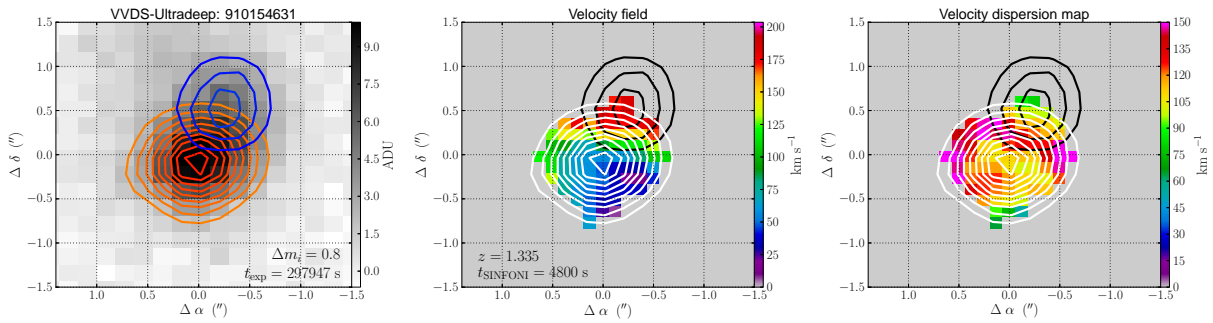


Fig. 18. The same as Fig. 3, but for the MASSIV source 910154631 (major merger). The outer contour marks the 1.9 ADU ($6\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 0.8 ADU ($2.5\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

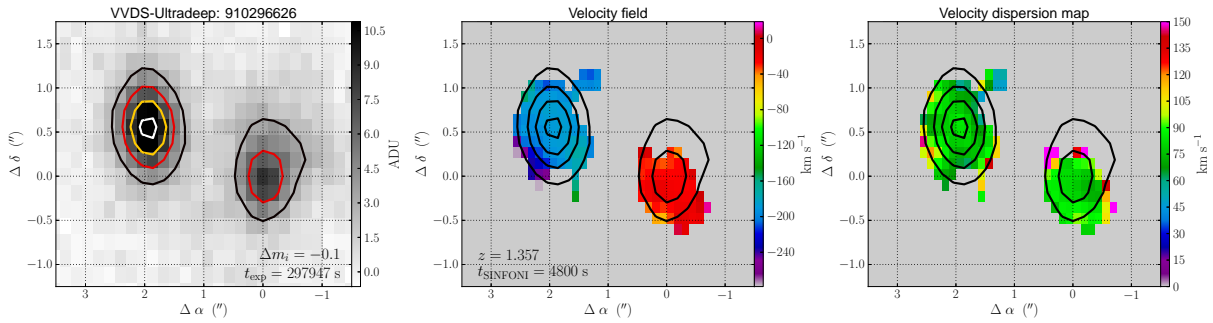


Fig. 19. The same as Fig. 8, but for the MASSIV source 910296626 (major merger). The outer contour marks the 3.32 ADU ($10\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 2.56 ADU ($8\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

GALFIT model with two components reproduces the shape of the velocity map and suggests that the extended region is due to the companion galaxy. The luminosity difference from the models is $\Delta m_i = 0.7$. Thus, we classify this system as a major merger.

- 020283830 (Fig. 14). *No major merger.* The velocity map shows two different components with a separation of $r_p = 8.5h^{-1}$ kpc and a relative velocity of $\Delta v \sim 500 \text{ km s}^{-1}$. The companion only presents six detected pixels in the $H\alpha$ map, already suggesting a minor companion. We used GALFIT to model the system, finding $\Delta m_i = 1.9$. However, the model of

the companion provided by GALFIT is a point-like source, thus overestimating the luminosity of the extended companion. The measurement with SExtractor in the image with the principal galaxy subtracted suggests $\Delta m_i = 2.1$. Since the companion is fainter than the major merger limit, we do not classify the system as a major merger.

- 020465775 (Fig. 15). **Major merger.** The velocity map suggests two projected components. The GALFIT model with two components finds that the companion galaxy is toward the north-west of the principal galaxy, at $r_p = 3.6h^{-1}$ kpc and $\Delta v \sim 40 \text{ km s}^{-1}$. The $H\alpha$ emission of this system is located

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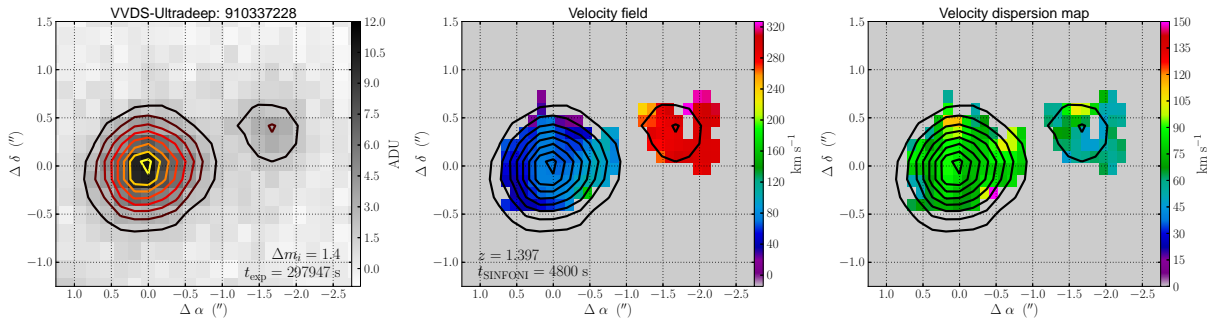


Fig. 20. The same as Fig. 8, but for the MASSIV source 910337228 (major merger). The outer contour marks the 1.73 ADU ($4\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 1.30 ADU ($3\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

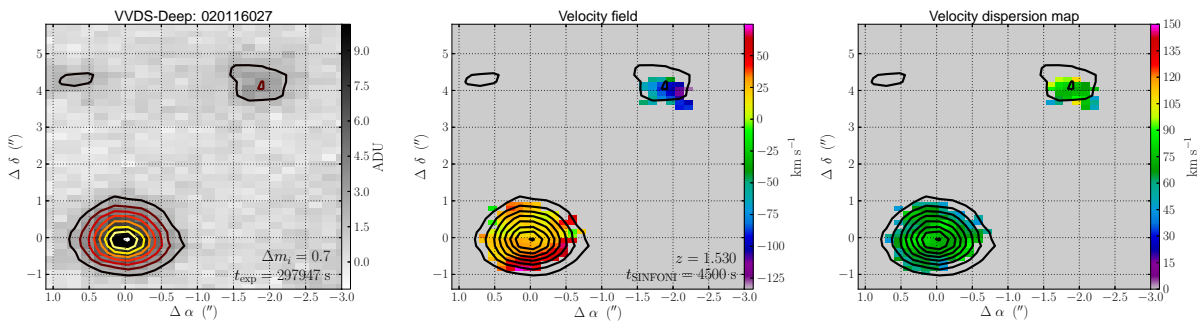


Fig. 21. The same as Fig. 8, but for the MASSIV source 020116027 (major merger). The principal galaxy is that in the south-east. The outer contour marks the 2.02 ADU ($3.5\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 1.16 ADU ($2\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

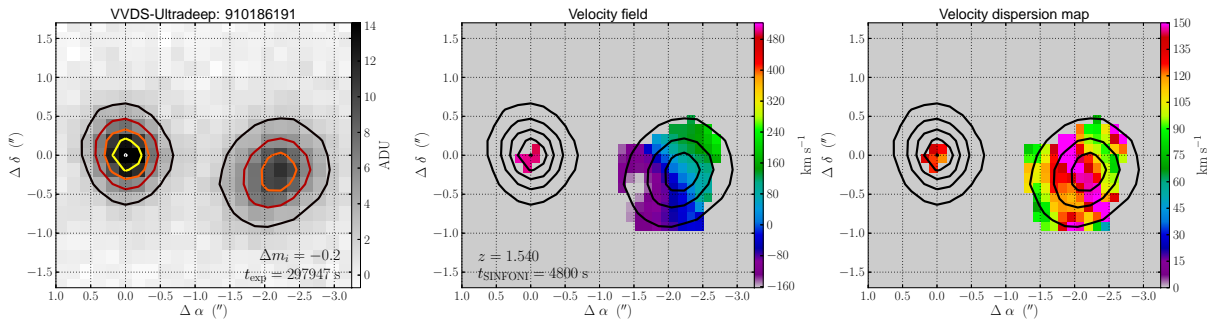


Fig. 22. The same as Fig. 3, but for the MASSIV source 910186191 (major merger). The outer contour marks the 2.83 ADU ($8\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 2.83 ADU ($8\sigma_{\text{sky}}$) steps. [A colour version of this plot is available at the electronic edition].

in the central part of the galaxies, since the i -band models match the velocity map at $9\sigma_{\text{sky}}$ level, i.e., we do not detect emission from the outer parts of the galaxies. The position of the companion explains the abnormal velocity pattern and the high velocity dispersion peak in the maps. The luminosity difference of the system is $\Delta m_i = 0.7$, so we classify it as a major merger.

- 220376206 (Fig. 16). *No major merger.* The velocity map shows two different components with a separation of $r_p = 13.4h^{-1}$ kpc and a relative velocity of $\Delta v \sim 400$ km s $^{-1}$. The companion, located toward the north, presents ten detected

pixels and its velocity is inconsistent with that expected from the velocity field of the principal galaxy. We find $\Delta m_i = 2.4$, so the system is not a major merger.

- 220544103 (Fig. 17). **Major merger.** The velocity map suggests two projected components. The southern component presents a large velocity gradient and defines the kinematical centre of the system, while the northern component is more extended and has a nearly flat velocity field. The GALFIT model with two components recovers the configuration in the velocity map and suggests that the southern component is edge-on, while the northern component is nearly face-on,

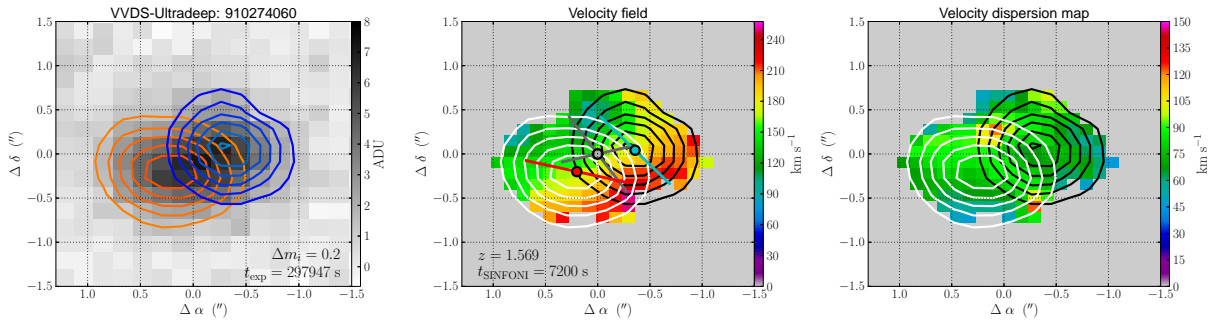
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Fig. 23. The same as Fig. 3, but for the MASSIV source 910274060 (major merger). The outer contour marks the 1.20 ADU ($3\sigma_{\text{sky}}$) isophote, while brighter isophotes increase in 1.00 ADU ($2.5\sigma_{\text{sky}}$) steps. The grey dot in the central figure marks the centre of the system given by the kinematical model, while the red/blue dot marks the photometric centre of the principal/companion galaxy given by the two-component model of the source. The long bars mark the position angle of the previous models. The short bar marks the photometric position angle given by the one-component model of the source. [A colour version of this plot is available at the electronic edition].

thus explaining the observed high and null velocity gradients. The separation between the components is $r_p = 5h^{-1}$ kpc, while their relative velocity is $\Delta v \sim 75 \text{ km s}^{-1}$. From the GALFIT models we estimate $\Delta m_i = -1.1$, and we classify the system as a major merger.

- 910154631 (Fig. 18). **Major merger.** The velocity map suggests two projected components, with the companion toward the north-west. The i -band image shows another two well separated sources close to the MASSIV target. These sources are not detected in $H\alpha$. To avoid contamination from these sources in the i -band photometry, we performed a four component fitting, with two components for the MASSIV target and one component for each nearby source. The GALFIT model finds the second component of the MASSIV target at the expected position, but we only detect the southern half of the i -band source in $H\alpha$. We explored the reduced data cube of the source, and we find (i) there are two clear velocity gradients in the cube, reinforcing the presence of two different components and (ii) there is an OH sky-line in the channels in which we expect the northern part of the companion, explaining the non detection in the maps. In addition, the higher velocity dispersion of the MASSIV target occurs in the expected overlapping region between both components. We conclude that this is a close pair system with $r_p = 4.2h^{-1}$ kpc, $\Delta v \sim 130 \text{ km s}^{-1}$ and $\Delta m_i = 0.8$. Thus, we classify the system as a major merger.
- 910296626 (Fig. 19). **Major merger.** The velocity map shows two different components with a separation of $r_p = 12.1h^{-1}$ kpc and a relative velocity of $\Delta v \sim 165 \text{ km s}^{-1}$. The companion is located toward the north-east. We find $\Delta m_i = -0.1$ and $\Delta m_{K_s} = -0.2$, so the system is a major merger.
- 910337228 (Fig. 20). **Major merger.** The velocity map shows two different components with a separation of $r_p = 9.5h^{-1}$ kpc and a relative velocity of $\Delta v \sim 220 \text{ km s}^{-1}$. The companion is toward the west and has $\Delta m_i = 1.4$. Thus, we classify the system as a major merger.

In summary for this redshift range, we classify 7 of the 11 close pair candidates as major mergers. This translates to a gas-rich major merger fraction of $f_{\text{MM}} = 0.201^{+0.080}_{-0.051}$ at $\bar{z}_{r,2} = 1.32$.

4.3. Close pair candidates at $1.5 \leq z < 1.8$

The weighted mean redshift of the third redshift bin is $\bar{z}_{r,3} = 1.54$. This is a redshift range where there is no measurement of the merger fraction from spectroscopic close pairs yet. We identify 3 close pair candidates over 12 galaxies:

- 020116027 (Fig. 21). **Major merger.** The velocity map shows two different components with a separation of $r_p = 26.8h^{-1}$ kpc and a relative velocity of $\Delta v \sim 100 \text{ km s}^{-1}$. The companion is toward the north-west. The luminosity difference is $\Delta m_i = 0.7$, suggesting a major merger. The difference in the K_s band is $\Delta m_{K_s} = 0.5$, confirming the previous major merger classification.
- 910186191 (Fig. 22). **Major merger.** The velocity map shows two different components with a separation of $r_p = 12.7h^{-1}$ kpc and a relative velocity of $\Delta v \sim 450 \text{ km s}^{-1}$. The MASSIV target is only detected in 6 pixels because there is an OH sky-line in the position of $H\alpha$ at its redshift. The companion, located toward the west, is well detected. We find $\Delta m_i = -0.2$, this is, the companion is slightly brighter than the principal galaxy. However, the MASSIV target is barely detected in the K_s band, with $\Delta m_{K_s} = -2.4$, suggesting a low mass system. This is the only system in which the classification in the two bands is different. Fortunately, both sources are VVDS targets, and we have an estimation of their stellar masses from SED fitting. The difference in stellar mass is $\mu \sim 1/3$, so we classify the system as a major merger.
- 910274060 (Fig. 23). **Major merger.** The velocity map is consistent with one single component. However, the position angle (PA) from the i -band photometry, $\text{PA} = 105^\circ$ (North has $\text{PA} = 0^\circ$ and East has $\text{PA} = 90^\circ$), is nearly perpendicular to that from the kinematical modelling, $\text{PA} = 33^\circ$. This suggests a complex system, so we performed the GALFIT modelling with two sources. We recover well two sources, one in the north and the other in the south. The photometric PAs of these two sources, provided by the GALFIT fitting, are now in better agreement with the kinematical one ($\text{PA}_1 = 75^\circ$, $\text{PA}_2 = 41^\circ$), supporting that this is a close pair system. The separation between the components is $3.4h^{-1}$ kpc, with $\Delta v \sim 10 \text{ km s}^{-1}$, and the luminosity difference is $\Delta m_i = 0.2$. Thus, we classify the system as a major merger.

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In summary, we identify the 3 candidates as major mergers. This translates to a gas-rich major merger fraction of $f_{\text{MM}} = 0.323^{+0.201}_{-0.107}$. Note that in this range our merger candidates have $r_{\text{p}}^{\text{max}} = 30h^{-1}$ kpc to improve the statistics. Applying Eq. (7) we estimate $f_{\text{MM}} = 0.220^{+0.137}_{-0.073}$ for $r_{\text{p}}^{\text{max}} = 20h^{-1}$ kpc at $\bar{z}_{1,3} = 1.54$.

Our results alone, summarised in Table 2, suggest a constant major merger fraction of $f_{\text{MM}} \sim 0.21$ at $0.9 < z < 1.8$ for $r_{\text{p}}^{\text{max}} = 20h^{-1}$ kpc close pairs (Fig. 24). *This merger fraction at $z > 1$ is higher by an order of magnitude than in the local universe, where $f_{\text{MM}} \sim 0.01 - 0.03$ (Patton et al. 2000; De Propriis et al. 2007; Patton & Atfield 2008; Domingue et al. 2009; Darg et al. 2010; Xu et al. 2012). This is the first main result of the present paper. We compare our major merger fractions with others in the literature in Sect. 6.1.*

5. The gas-rich major merger rate in MASSIV

In this section we estimate the gas-rich major merger rate (R_{MM}), defined as the number of mergers per galaxy and Gyr, of star-forming galaxies at $0.9 < z < 1.8$. We remind here the steps to transform a merger fraction to a merger rate. Following de Ravel et al. (2009), we define the major merger rate as

$$R_{\text{MM}} = C_{\text{m}} f_{\text{MM}} T_{\text{MM}}^{-1}, \quad (6)$$

where the factor C_{m} is the fraction of the observed close pairs that finally merge in a typical time scale T_{MM} . The typical merger time scale can be estimated by cosmological and N -body simulations. In our case, we compute the major merger time scale from the cosmological simulations of Kitzbichler & White (2008), based on the Millennium simulation (Springel et al. 2005). This major merger time scale refers to major mergers ($\mu \geq 1/4$ in stellar mass), and depends mainly on $r_{\text{p}}^{\text{max}}$ and on the stellar mass of the principal galaxy, with a weak dependence on redshift in our range of interest (see de Ravel et al. 2009, for details). We measured the median-weighted stellar mass from MASSIV sources in each of the three redshift bins under study, and estimated the merger time scale for these stellar masses. These time scales already include the factor C_{m} (see Patton & Atfield 2008; Bundy et al. 2009; Lin et al. 2010; López-Sanjuan et al. 2011), so we take $C_{\text{m}} = 1$ in the following. In addition, López-Sanjuan et al. (2011) show that the time scales from Kitzbichler & White (2008) are equivalent to that from the N -body/hydrodynamical simulations by Lotz et al. (2008). However, we stress that these merger time scales have an additional factor of two uncertainty in their normalization (e.g., Hopkins et al. 2010; Lotz et al. 2011). We summarise the stellar masses, the merger time scales and the gas-rich major merger rates in Table 2. As for the merger fraction, MASSIV data suggests a nearly constant major merger rate at $0.9 < z < 1.8$, $R_{\text{MM}} \sim 0.12 \text{ Gyr}^{-1}$ (Fig. 25). We study in detail the evolution of the major merger rate at $z \lesssim 1.5$ in Sect. 6.2.

6. The redshift evolution of the gas-rich major merger fraction and rate up to $z \sim 1.5$

In this section we use the MASSIV results at $z > 1$ to expand the study of the gas-rich major merger fraction (Sect. 6.1) and rate (Sect. 6.2) from spectroscopic close pairs to the redshift desert. Then, we explore the importance of gas-rich major mergers in the assembly of the red sequence since $z \sim 1.5$ in Sects. 6.3 and 6.4.

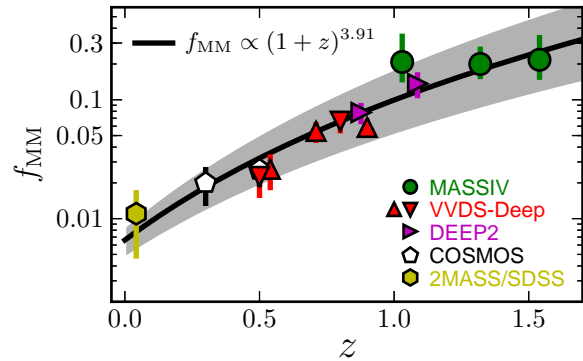


Fig. 24. Gas-rich major merger fraction of $\bar{M}_{\star} \sim 10^{10-10.5} M_{\odot}$ galaxies as a function of redshift. Circles are from this MASSIV data set, triangles are from de Ravel et al. (2009) and inverted triangles are from López-Sanjuan et al. (2011), both in VVDS-Deep, right-pointing triangles are from Lin et al. (2008) in DEEP2 redshift survey, pentagons are from Xu et al. (2012) in the COSMOS field, and the hexagon is from Xu et al. (2012) in 2MASS/SDSS. The solid line is the least-squares fit of a power-law function, $f_{\text{MM}} = 0.0066 \times (1+z)^{3.91}$, to the data. The grey area marks the 3σ confidence interval in the fit. [A colour version of this plot is available at the electronic edition].

6.1. The redshift evolution of the gas-rich major merger fraction

In this section we compare the merger fraction from MASSIV with those from previous works. Because the merger fraction evolution depends on mass (e.g., de Ravel et al. 2009, 2011), luminosity (e.g., de Ravel et al. 2009; López-Sanjuan et al. 2010a) and colour (e.g., Lin et al. 2008; Chou et al. 2011), we focus on samples with $\bar{M}_{\star} \sim 10^{10-10.5} M_{\odot}$ (Salpeter 1955 IMF) to minimise systematics. In addition, this mass regime is greatly dominated by gas-rich (wet) mergers, as those that we observe in MASSIV, at least at $z \gtrsim 0.2$ (Lin et al. 2008; de Ravel et al. 2009; Chou et al. 2011).

We define the major ($\mu \geq 1/4$) merger fraction normalised to $r_{\text{p}}^{\text{max}} = 20h^{-1}$ kpc as

$$f_{\text{MM}}(20, 1/4) = C_{\text{p}} F(\mu) \left(\frac{20h^{-1} \text{ kpc}}{r_{\text{p}}^{\text{max}}} \right)^{0.95} f_{\text{m}}(r_{\text{p}}^{\text{max}}, \mu), \quad (7)$$

where the factor $C_{\text{p}} = r_{\text{p}}^{\text{max}} / (r_{\text{p}}^{\text{max}} - r_{\text{p}}^{\text{min}})$ accounts for the missing close companions at small radii in those studies with $r_{\text{p}}^{\text{min}} > 0$ (e.g., Bell et al. 2006) and the factor $F(\mu)$ translates the merger fraction for a given μ to the major merger fraction. The merger fraction depends on μ as $f_{\text{m}}(\geq \mu) \propto \mu^s$ (e.g., López-Sanjuan et al. 2011), that implies $F(\mu) = (4\mu)^{-s}$. We take $s = -0.9 \pm 0.4$, a value derived from the observational estimations of López-Sanjuan et al. (2011, 2012) and Xu et al. (2012). The search radius dependence of the major merger fraction, $f_{\text{MM}} \propto r_{\text{p}}^{0.95}$, is the observational one found by López-Sanjuan et al. (2011) in the VVDS. With Eq. (7) we avoid systematic differences due to the close pair definition when comparing different works.

de Ravel et al. (2009) study the major merger fraction of $M_{\star} \geq 10^{9.75} M_{\odot}$ ($\bar{M}_{\star} \sim 10^{10.25} M_{\odot}$) galaxies in VVDS-Deep by spectroscopic close pairs, while López-Sanjuan et al. (2011) provide the major merger fraction of blue (star-forming)

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Table 2. Gas-rich major merger fraction and rate of star-forming galaxies at $0.9 < z < 1.8$ in the MASSIV sample.

z_r	N	N_p	\bar{z}_r	$\log(\bar{M}_*/M_\odot)$	r_p^{\max} (h^{-1} kpc)	T_{MM} (Gyr)	f_{MM}	R_{MM} (Gyr^{-1})
$0.94 \leq z \leq 1.06$	18	3	1.03	10.17	20	1.80	$0.208^{+0.152}_{-0.068}$	$0.116^{+0.084}_{-0.038}$
$1.2 \leq z < 1.5$	30	7	1.32	10.57	20	1.37	$0.201^{+0.080}_{-0.051}$	$0.147^{+0.058}_{-0.037}$
$1.5 \leq z < 1.8$	12	3	1.54	10.09	30	2.54	$0.323^{+0.201}_{-0.107}$	$0.127^{+0.079}_{-0.042}$

galaxies with $\bar{M}_* \sim 10^{10.55} M_\odot$ in the same sample. In both studies $r_p^{\max} = 100h^{-1}$ kpc. Lin et al. (2008) report the number of companions of $-21 \leq M_B + 1.3z \leq -19$ galaxies ($\bar{M}_* \sim 10^{10.25} M_\odot$) with $10h^{-1}$ kpc $\leq r_p \leq 30h^{-1}$ kpc in three DEEP2 redshift survey (Newman et al. 2012b) fields. Their principal and companion sample are the same, so they miss major companions near to the selection boundary. Thus, we apply an extra factor 1.74 to Eq. (7) to account for these missing companions (see Lin et al. 2004, for details). Xu et al. (2012) measure the fraction of galaxies in close pairs with $\mu \geq 1/2.5$ in the COSMOS⁴ (Cosmological Evolution Survey, Scoville et al. 2007) and SDSS⁵ (Sloan Digital Sky Survey, Abazajian et al. 2009) surveys for $\bar{M}_* \sim 10^{10.2} M_\odot$ galaxies. We applied a factor 0.5 to pass from their number of galaxies in close pairs to the number of close pair systems in the sample (C. K. Xu, private communication), and a factor $F(1/2.5) = 1.5 \pm 0.3$ to obtain the major merger fraction. All these published (gas-rich) major merger fractions are shown as a function of redshift in Fig. (24), together with the values derived from MASSIV.

We parametrise the redshift evolution of the (gas-rich) major merger fraction with a power-law,

$$f_{MM} = f_{MM,0} (1+z)^m. \quad (8)$$

The least-squares fit to all the data in Fig. 24 yields $f_{MM,0} = (6.6 \pm 0.6) \times 10^{-3}$ and $m = 3.91 \pm 0.16$. We find good agreement between all works, with the MASSIV point at $z \sim 1$ being higher than expected from the fit, but consistent within errors with the measurement of Lin et al. (2008) at that redshift. In the next section we show that this difference disappears when the stellar mass of the samples is taken into account, emphasizing the importance of comparing results from similar parent samples.

6.2. The redshift evolution of the gas-rich major merger rate

We use Eq. (6) to translate the original (i.e., without any normalization in r_p^{\max}) major merger fractions reported in previous section into merger rates. We show them in Fig. 25. The good agreement between different works is remarkable, reinforcing the idea that the merger time scales used account properly for the dependence of the merger fraction both on r_p^{\max} and on stellar mass.

We also show the major merger rate from morphological criteria derived by López-Sanjuan et al. (2009b). They measure the gas-rich merger fraction of $\bar{M}_* \sim 10^{10.5} M_\odot$ galaxies from asymmetries (A) in the GOODS⁶ (Giavalisco et al. 2004) South field and take into account the effect of observational

⁴ <http://cosmos.astro.caltech.edu>

⁵ <http://www.sdss.org/>

⁶ <http://www.stsci.edu/science/goods/>

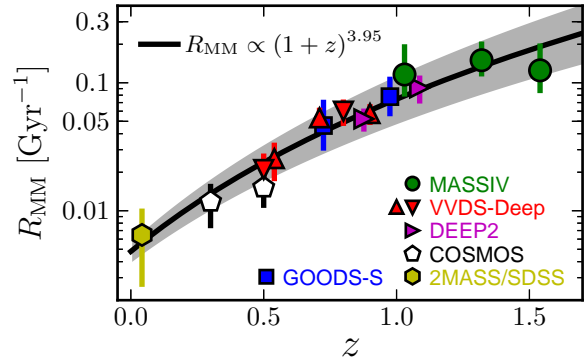


Fig. 25. Gas-rich major merger rate of $\bar{M}_* \sim 10^{10-10.5} M_\odot$ galaxies as a function of redshift. Circles are from this MASSIV data set, triangles are from de Ravel et al. (2009) and inverted triangles are from López-Sanjuan et al. (2011), both in VVDS-Deep, right-pointing triangles are from Lin et al. (2008) in DEEP2 redshift survey, squares are from López-Sanjuan et al. (2009b) in GOODS-S from morphological criteria, pentagons are from Xu et al. (2012) in the COSMOS field, and the hexagon is from Xu et al. (2012) in 2MASS/SDSS. The solid line is the least-squares fit of a power-law function, $R_{MM} = 0.0048 \times (1+z)^{3.95}$, to the data. The grey area marks the 3σ confidence interval in the fit. [A colour version of this plot is available at the electronic edition].

errors in z and A , that lead to overestimations in the major merger fraction by a factor of two-three (López-Sanjuan et al. 2009a,b), using maximum likelihood techniques developed in López-Sanjuan et al. (2008). Other studies find good agreement between the asymmetry-based major merger rates from López-Sanjuan et al. (2009b) and those from close pair statistics (López-Sanjuan et al. 2010a; Lotz et al. 2011; de Ravel et al. 2011), confirming the robustness of their methodology.

We also parametrise the redshift evolution of the (gas-rich) major merger rate with a power-law,

$$R_{MM} = R_{MM,0} (1+z)^n. \quad (9)$$

The least-squares fit to all the data in Fig. 25 yields $R_{MM,0} = (4.8 \pm 0.3) \times 10^{-3} \text{ Gyr}^{-1}$ and $n = 3.95 \pm 0.12$. The agreement between different works points out the importance of comparing results from similar parent samples to avoid systematics (see also Lotz et al. 2011, for an extensive discussion on this topic). We also point out that the merger rate and the merger fraction of star-forming galaxies show a similar evolution with redshift.

The second main result in this paper is that *the major merger rate is well described by a power-law function up to*

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$z \sim 1.5$. However, we note that our MASSIV data seem to indicate a flattening of the merger rate's evolution beyond $z \sim 1$. Previous studies from morphological criteria (e.g., Conselice 2003; Conselice et al. 2008) and from photometric pairs (Ryan et al. 2008) suggest that the power-law parametrization is not longer valid at $z \gtrsim 1.5$, where a lower merger fraction than expected from the low- z evolution is measured, possibly indicating a maximum in the major merger rate at $z \sim 2$ (e.g., Conselice 2006; Ryan et al. 2008; López-Sanjuan et al. 2009b). Our new measurements agree with this picture and measurements from spectroscopic close pairs beyond $z \sim 1.5$ are needed to test the early evolution of the merger fraction.

The power-law index $n = 3.95 \pm 0.12$ is higher than several previous measurements in the literature (e.g., Bridge et al. 2010; Lotz et al. 2011), as well as our major merger fraction evolution, $m = 3.91 \pm 0.16$. However, our results refer to $\bar{M}_* \sim 10^{10-10.5} M_\odot$ star-forming galaxies, and it is known that the merger fraction and rate evolve faster for blue, star-forming galaxies than for the red and global populations (e.g., Lin et al. 2008; de Ravel et al. 2009; Chou et al. 2011; López-Sanjuan et al. 2011).

6.3. Number of gas-rich mergers since $z = 1.5$

We can obtain the average number of gas-rich major mergers per star-forming galaxy between z_2 and $z_1 < z_2$ as

$$N_{\text{MM}}(z_1, z_2) = \int_{z_1}^{z_2} \frac{R_{\text{MM}} dz}{(1+z)H_0 E(z)}, \quad (10)$$

where $E(z) = \sqrt{\Omega_\Lambda + \Omega_m(1+z)^3}$ in a flat universe. Using the merger rate parametrisation in Eq. (9), we obtain $N_{\text{MM}}(0, 1.5) = 0.35 \pm 0.04$. Interestingly, half of this merging activity happens at $z > 1$, with $N_{\text{MM}}(1, 1.5) = 0.18 \pm 0.02$ and $N_{\text{MM}}(0, 1) = 0.17 \pm 0.02$. Because the cosmic time lapse in these redshift intervals is 1.55 Gyr and 7.7 Gyr, respectively, the average merger activity was higher at $1 < z < 1.5$ than at $1 < z$ by a factor of five. In the next section we further explore the consequences of this very different major merger activity above and below $z \sim 1$ for the assembly of the red sequence.

6.4. Testing the major merger origin of massive early-type galaxies

The number density of massive ($M_* \gtrsim 10^{11} M_\odot$) early-type galaxies (E/S0, ETGs in the following) has increased with cosmic time since $z \sim 3$ (e.g., Pozzetti et al. 2010; Buitrago et al. 2013), with ETGs being the dominant population among massive galaxies only since $z \sim 1$ (Vergani et al. 2008; Buitrago et al. 2013; van der Wel et al. 2011; van Dokkum et al. 2011). Gas-rich major mergers have been proposed as an efficient mechanism to transform star-forming late-type galaxies into red ETGs (e.g., Naab et al. 2006; Rothberg & Joseph 2006a,b; Hopkins et al. 2008; Rothberg & Fischer 2010; Bournaud et al. 2011), so the comparison between the observed number density evolution of ETGs (ρ_{ETG}) and the major merger history of star-forming galaxies imposes important constraints on the role of mergers in galaxy evolution.

In their work, Robaina et al. (2010) and Man et al. (2012) integrate the observed major merger rate over cosmic time and predict the evolution of the number density of massive galaxies assuming that all the merger remnants are new massive galaxies. They suggest that major mergers are common enough to explain the number density evolution of massive galaxies (ETGs

+ spirals) at $z < 1$ and $z < 3$, respectively. Eliche-Moral et al. (2010) model the evolution of the luminosity function backwards in time for bright galaxies, selected according to their colours (red/blue/total) and their morphologies. They find that the observed luminosity functions' evolution can be naturally explained by the observed gas-rich and dry major merger rates, and that 50–60% of massive ETGs in the local universe were formed by major mergers at $0.8 < z < 1$, with a small number evolution since $z = 0.8$.

In this section we implement a model to explore the role of mergers in the number density evolution of massive ETGs since $z \sim 1.3$. As reference values, we use the number densities of massive ($M_* \geq 10^{11.25} M_\odot$) ETGs provided by Buitrago et al. (2013). They perform a consistent morphological study by visual inspection between $z = 0$ and $z = 3$, combining the SDSS, POWIR (Palomar Observatory Wide-Field Infrared, Conselice et al. 2007) and GNS⁷ (GOODS NICMOS Survey, Conselice et al. 2011a) surveys. We also use the number densities of massive spheroidal galaxies with $M_* \geq 10^{11.25} M_\odot$, selected by automatic indices in the zCOSMOS⁸ (Lilly et al. 2007) survey, provided by Pozzetti et al. (2010). We show these number densities in Fig. 26.

In our model we assume that, after the final coalescence of the merging galaxies, a cosmic time $\Delta t = 0.5$ Gyr is necessary for the merger remnant to be classified as an ETG (see Eliche-Moral et al. 2010, for a detailed summary of this topic). This implies that the new ETGs which appeared between z_{max} and z_{min} came from the merger activity in the range $z \in [z_1, z_2]$, where $z_1 = z_{\text{min}} + \Delta z$, $z_2 = z_{\text{max}} + \Delta z$, and Δz is the redshift interval that spans Δt in each case. Therefore, we take $z_{\text{max}} = 1.3$ as the upper redshift in our model because that implies $z_2 \sim 1.5$, the redshift limit of the present merger rate study (see Sect. 6.2).

The number density of new ETGs with stellar mass $M_* \geq M_{*,\text{lim}}$ from gas-rich major merger events appeared in the range $z_{\text{min}} \leq z < z_{\text{max}}$ is

$$\rho_{\text{wet}}(z_{\text{min}}, z_{\text{max}}, M_{*,\text{lim}}) = \int_{z_1}^{z_2} \int_0^\infty \Phi R_{\text{MM}} E_{\text{MM}} f_{\text{LTG}} dM_* dz, \quad (11)$$

where the different elements in the integral are (Fig. 27):

- the global mass function, parametrised with a Schechter function

$$\Phi(z, M_*) = \frac{\phi^*(z)}{M_*^*(z)} \left(\frac{M_*}{M_*^*(z)} \right)^{\alpha(z)} \exp\left(-\frac{M_*}{M_*^*(z)}\right). \quad (12)$$

We assumed the Schechter function parameters from Pérez-González et al. (2008) and used their parametrisation with redshift provided by López-Sanjuan et al. (2010a),

$$\log(\phi^*(z)/\text{Mpc}^{-3}) = -2.72 - 0.56(z - 0.5), \quad (13)$$

$$\log(M_*^*(z)/M_\odot) = 11.23 + 0.13(z - 0.5), \quad (14)$$

$$\alpha(z) = -1.22 - 0.04(z - 0.5). \quad (15)$$

- The gas-rich major merger rate, $R_{\text{MM}}(z)$, as measured in Sect. 6.2. We assumed that it is independent of the stellar mass.
- The merger efficiency function, $E(z, M_*, M_{*,\text{lim}}, \mu_{\text{max}}, \mu_{\text{min}})$. This function provides the probability that a gas-rich merger with a mass ratio $\mu_{\text{min}} < \mu \leq \mu_{\text{max}}$ produces an early-type remnant more massive than $M_{*,\text{lim}}$. The merger efficiency function takes several effects into account:

⁷ <http://www.nottingham.ac.uk/astronomy/gns/>

⁸ <http://www.astro.phys.ethz.ch/zCOSMOS/>

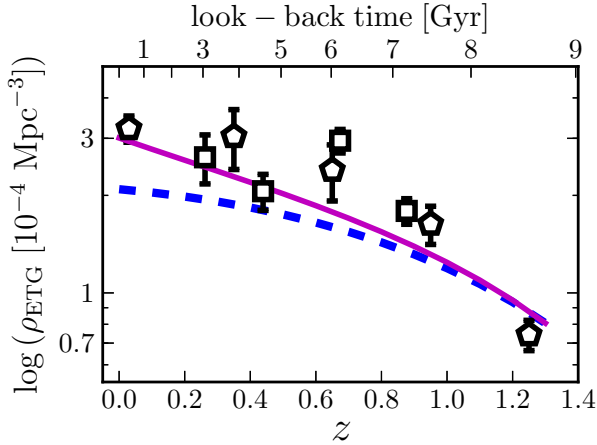


Fig. 26. Number density evolution of massive ($M_* \geq 10^{11.25} M_\odot$) ETGs (E/S0) as a function of redshift from Buitrago et al. (2013, pentagons) and Pozzetti et al. (2010, squares). The dashed line is the expected number density evolution due to gas-rich (wet) major mergers from our model. The solid line is the expected number density evolution due to wet major and dry mergers (both major and minor) from our model. Mergers are common enough to drive the number density evolution of massive ETGs since $z \sim 1.3$. [A colour version of this plot is available at the electronic edition].

- The stellar mass difference μ needed to reach $M_{*,\text{lim}}$ for a given M_* . For example, and regarding major mergers with $\mu \geq 1/4$, the $\mu_{\text{max}} = 1$ mergers of galaxies with $M_* \geq 10^{10.7} M_\odot$ will provide a massive remnant with $M_* \geq M_{*,\text{lim}} = 10^{11} M_\odot$, while for $\mu_{\text{min}} = 1/4$ only galaxies with $M_* \geq 10^{10.9} M_\odot$ can be progenitors of massive ETGs. Thus, mergers below $M_{*,\text{lim}}/(1 + \mu_{\text{min}})$ do not contribute to the assembly of massive ETG, while all mergers above $M_{*,\text{lim}}/(1 + \mu_{\text{max}})$ contribute. Between μ_{min} and μ_{max} , we used the dependence of the merger fraction on the mass ratio μ , $f_{\text{MM}} \propto \mu^s$ (e.g., López-Sanjuan et al. 2011), to estimate the effective merger rate at $M_* = M_{*,\text{lim}}/(1 + \mu)$.
- The extra stellar mass in the final remnant due to the star formation during the merger, f_{sf} . This extra stellar mass simply decreases the limiting mass for a given μ described in the previous item by a factor $(1 + f_{\text{sf}})^{-1}$.
- The fraction of remnants of a gas-rich major merger that are ETGs. We express this fraction as $1 - f_{\text{disc}}$, where f_{disc} is the fraction of remnants that rebuild a disc component and are classified as late-type galaxies after the merger (see Sect. 6.4.2, for further details).

Formally, the merger efficiency function is defined as

$$E = \begin{cases} 1 - f_{\text{disc}}, & \text{if } M_* \geq M_{*,1}, \\ (1 - f_{\text{disc}}) \times (\mu/\mu_{\text{min}})^s, & \text{if } M_{*,2} \leq M_* < M_{*,1}, \\ 0, & \text{if } M_* < M_{*,2}, \end{cases} \quad (16)$$

where

$$M_{*,1} = \frac{M_{*,\text{lim}}}{(1 + \mu_{\text{min}})(1 + f_{\text{sf}})}, \quad (17)$$

$$M_{*,2} = \frac{M_{*,\text{lim}}}{(1 + \mu_{\text{max}})(1 + f_{\text{sf}})}, \quad (18)$$

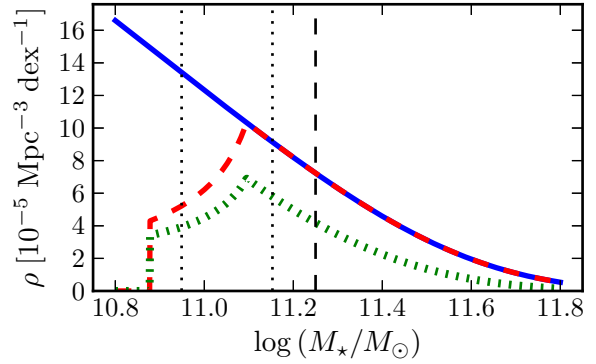


Fig. 27. Number density distribution from the different functions in Eq. (11) at $z = 1$. The solid line is the stellar mass function (Φ) multiplied by the major merger rate (R_{MM}). The dashed line includes the major merger efficiency function (E_{MM}). The dotted line includes the fraction of late-type galaxies (f_{LTG}) and provides the final function that we integrate over stellar mass and cosmic time to obtain ρ_{wet} . The vertical dashed line marks the limiting stellar mass that define massive galaxies, $M_{*,\text{lim}} = 10^{11.25} M_\odot$, while the dotted lines mark the stellar masses $M_{*,1} = 10^{10.15} M_\odot$ and $M_{*,2} = 10^{11.05} M_\odot$ when $f_{\text{sf}} = 0$ (see text for details). [A colour version of this plot is available at the electronic edition].

$$\mu = \frac{M_{*,\text{lim}}}{M_*(1 + f_{\text{sf}})} - 1. \quad (19)$$

For convenience, we define the major merger efficiency in Eq. (11) as $E_{\text{MM}} \equiv E(z, M_*, M_{*,\text{lim}}, \mu_{\text{max}} = 1, \mu_{\text{min}} = 1/4)$. We assumed that $f_{\text{disc}} = 0$, i.e., all the merger remnants are ETGs (see Sect. 6.4.2), and that $s = -0.9$ as in Sect 6.1. To estimate f_{sf} we used the parametrisation of the gas fraction as a function of stellar mass and cosmic time provided by Rodrigues et al. (2012),

$$f_{\text{gas}}(z, M_*) = \frac{M_{\text{gas}}}{M_* + M_{\text{gas}}} = \frac{1}{1 + [M_*/10^{A(t)}]^{B(t)}}, \quad (20)$$

where t is the cosmic time between redshift z and the present,

$$A(t) = 9.15 + 0.13t, \quad (21)$$

$$B(t) = 0.5 + 13.36 \times \exp(-38.02/t). \quad (22)$$

$A(t)$ represents the stellar mass at a given time for which the gas fraction is equal to 50%. The parameter increases linearly with lookback time. $B(t)$ corresponds to the slope of the function. Then, we assumed the prescriptions in Hopkins et al. (2009) to estimate the amount of the initial gas mass that is transformed into stars during the merger, $f_{\text{burst}} = f_{\text{gas}}(1 - f_{\text{gas}})2\mu/(1 + \mu)$. Finally,

$$f_{\text{sf}} = M_{\text{burst}}/M_* = \frac{2}{3}f_{\text{gas}}, \quad (23)$$

where the factor $2/3$ is the efficiency for $\mu = 1/2$ mergers, which is the typical major merger mass ratio (López-Sanjuan et al. 2011, 2012).

- The fraction of late-type galaxies (spirals and irregulars), $f_{\text{LTG}}(z, M_*)$, is the number of potential gas-rich merger progenitors over the total population. We parametrise this fraction as

$$f_{\text{LTG}}(z, M_*) = 0.56 + 0.16z - 0.58[\log(M_*/M_\odot) - 11]. \quad (24)$$

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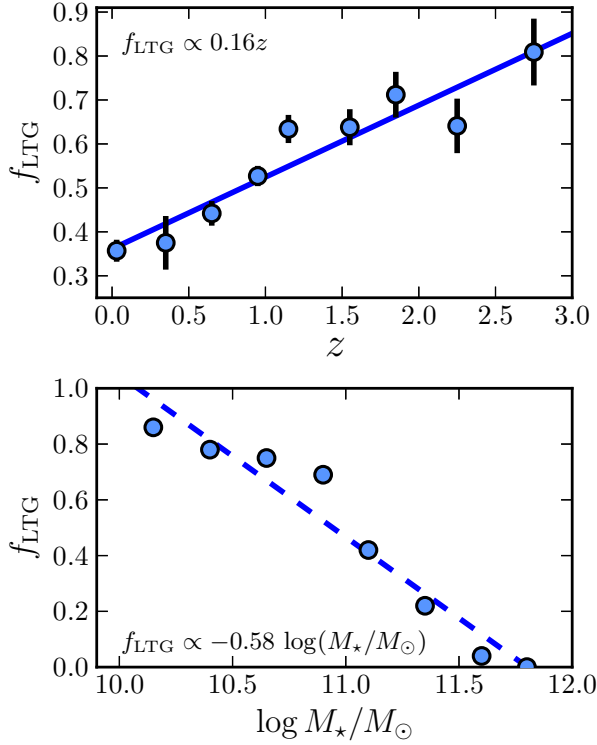


Fig. 28. Fraction of late-type galaxies as a function of redshift (top panel) and stellar mass at $z \sim 0$ (bottom panel). The redshift data points are from Buitrago et al. (2013), and the stellar mass ones from Bernardi et al. (2010) in the SDSS. The line in both panels is the best least-squares linear fit to the data. [A colour version of this plot is available at the electronic edition].

We estimated the redshift dependence from the f_{LTG} reported by Buitrago et al. (2013) at $z < 3$ (Fig. 28, top panel), and the mass dependence from the late-type fractions in the SDSS presented in Bernardi et al. (2010, Fig. 28, bottom panel). We assumed $f_{\text{LTG}} = 1$ when Eq. (24) is higher than one, and $f_{\text{LTG}} = 0$ when it is negative.

Finally, we define the fraction of new massive ETGs due to gas-rich (wet) mergers as

$$f_{\text{wet}}(z_{\text{min}}, z_{\text{max}}) = \frac{\rho_{\text{wet}}(z_{\text{min}}, z_{\text{max}}, 10^{11.25} M_{\odot})}{\rho_{\text{ETG}}(z_{\text{min}}) - \rho_{\text{ETG}}(z_{\text{max}})}. \quad (25)$$

Our merger model finds $f_{\text{wet}}(0, 1.3) \sim 50\%$, while this fraction increases up to $f_{\text{wet}}(z, 1.3) \gtrsim 90\%$ when we focus on the high- z regime, $z \gtrsim 0.8$ (dashed line in Fig. 26). This indicates that gas-rich major mergers are common enough at $z \gtrsim 0.8$ to explain the observed increase in the number density of massive ETGs. However, at $z \lesssim 0.8$ only one third of the evolution can be accounted by these mergers, $f_{\text{wet}}(0, 0.8) \sim 40\%$. That supports the idea that gas-rich major merging is the main process involved in the assembly of the red sequence at $z \gtrsim 1$ (e.g., Ilbert et al. 2010; López-Sanjuan et al. 2010b; Eliche-Moral et al. 2010; Prieto et al. 2013).

In addition to wet mergers, dry mergers can also increase the number density of massive ETGs. In this case they promote ETGs with $M_{\star} < 10^{11.25} M_{\odot}$ to the massive regime. We estimate the contribution of both major and minor dry mergers as

$$\rho_{\text{dry}}^{+}(z_{\text{min}}, z_{\text{max}}, M_{\star, \text{lim}}) = \int_{z_1}^{z_2} \int_0^{M_{\star, \text{lim}}} \Phi R_{\text{MM}}^{\text{ETG}} E_{\text{MM}} f_{\text{ETG}} dM_{\star} dz + \int_{z_1}^{z_2} \int_0^{M_{\star, \text{lim}}} \Phi R_{\text{mm}}^{\text{ETG}} E_{\text{mm}} f_{\text{ETG}} dM_{\star} dz, \quad (26)$$

where $R_{\text{MM}}^{\text{ETG}}$ ($R_{\text{mm}}^{\text{ETG}}$) is the major (minor) merger rate of massive ETGs galaxies from López-Sanjuan et al. (2012), $f_{\text{ETG}} = 1 - f_{\text{LTG}}$, the minor merger efficiency function is defined as $E_{\text{mm}} \equiv E(z, M_{\star}, M_{\star, \text{lim}}, \mu_{\text{max}} = 1/4, \mu_{\text{min}} = 1/10)$, and we assumed $f_{\text{sf}} = 0$ in both major and minor merger efficiency functions. Note that the integration in mass space only reach $M_{\star, \text{lim}}$ and that the merger rates from López-Sanjuan et al. (2012) include also mixed mergers (ETGs - LTGs). However, dry mergers between two already massive ETGs decrease the number density ρ_{ETG} . We take this into account with the following function,

$$\rho_{\text{dry}}^{-}(z_{\text{min}}, z_{\text{max}}, M_{\star, \text{lim}}) = \int_{z_1}^{z_2} \int_{M_{\star, \text{lim}}}^{\infty} \Phi R_{\text{MM}}^{\text{ETG}} \epsilon_{\text{MM}} f_{\text{ETG}} dM_{\star} dz + \int_{z_1}^{z_2} \int_{M_{\star, \text{lim}}}^{\infty} \Phi R_{\text{mm}}^{\text{ETG}} \epsilon_{\text{mm}} f_{\text{ETG}} dM_{\star} dz, \quad (27)$$

where in this case the merger efficiency function has the form

$$\epsilon = \begin{cases} 0.65, & \text{if } M_{\star} \geq M_{\star, 3}, \\ 0.65 \times (\mu/\mu_{\text{min}})^{\text{s}}, & \text{if } M_{\star, 4} \leq M_{\star} < M_{\star, 3}, \\ 0, & \text{if } M_{\star} < M_{\star, 4}, \end{cases} \quad (28)$$

where

$$M_{\star, 3} = \mu_{\text{min}}^{-1} M_{\star, \text{lim}}, \quad (29)$$

$$M_{\star, 4} = \mu_{\text{max}}^{-1} M_{\star, \text{lim}}, \quad (30)$$

$$\mu = \frac{M_{\star, \text{lim}}}{M_{\star}}. \quad (31)$$

and the factor 0.65 is the fraction of companions of massive galaxies that are already early-type/red galaxies (López-Sanjuan et al. 2012; Newman et al. 2012a). As previously, we define the major merger efficiency as $\epsilon_{\text{MM}} \equiv \epsilon(z, M_{\star}, M_{\star, \text{lim}}, \mu_{\text{max}} = 1, \mu_{\text{min}} = 1/4)$ and the minor merger efficiency as $\epsilon_{\text{mm}} \equiv \epsilon(z, M_{\star}, M_{\star, \text{lim}}, \mu_{\text{max}} = 1/4, \mu_{\text{min}} = 1/10)$.

Analogous to the wet merger case, we define

$$f_{\text{dry}}(z_{\text{min}}, z_{\text{max}}) = \frac{\rho_{\text{dry}}(z_{\text{min}}, z_{\text{max}}, 10^{11.25} M_{\odot})}{\rho_{\text{ETG}}(z_{\text{min}}) - \rho_{\text{ETG}}(z_{\text{max}})} = \frac{\rho_{\text{dry}}^{+}(z_{\text{min}}, z_{\text{max}}, 10^{11.25} M_{\odot}) - \rho_{\text{dry}}^{-}(z_{\text{min}}, z_{\text{max}}, 10^{11.25} M_{\odot})}{\rho_{\text{ETG}}(z_{\text{min}}) - \rho_{\text{ETG}}(z_{\text{max}})} \quad (32)$$

to explore the role of dry mergers in massive ETGs assembly since $z = 1.3$. We find $f_{\text{dry}}(0, 1.3) \sim 40\%$. As shown in Fig. 26, dry mergers are more important at recent cosmic times due to the increase in the number density of ETGs, in contrast with the diminishing importance of gas-rich mergers. For example, we have $f_{\text{dry}}(0.8, 1.3) \sim 15\%$, while $f_{\text{dry}}(0, 0.8) \sim 45\%$.

The combined effect of gas-rich and dry mergers, $f_{\text{tot}} = f_{\text{wet}} + f_{\text{dry}}$, is able to explain most of the evolution in ρ_{ETG} since $z = 1.3$, with $f_{\text{tot}}(0, 1.3) \sim 90\%$. Thus, our model suggests merging as the main process in the assembly of massive ETGs since $z = 1.3$. Two thirds of the number density evolution is due to gas-rich major mergers, while one third is coming from major and minor dry mergers.

The measurement of the merger rate at $z \gtrsim 1.5$ is needed to fully constraint the role of gas-rich major mergers in the early assembly of the red sequence, as well as the possible contribution of cold gas accretion in this mass assembly.

6.4.1. Fast and slow rotators in the local universe

The results from the SAURON project (de Zeeuw et al. 2002) propose a kinematical classification of morphological ETGs into fast and slow rotators (Emsellem et al. 2007). Recently, the ATLAS^{3D,9} (Cappellari et al. 2011) survey has observed a representative sample of 260 nearby ETGs, finding that the fraction of slow rotators increases with the dynamical mass (Emsellem et al. 2011). The cosmological simulation analysed by Khochfar et al. (2011) suggests that the main difference between fast and slow rotators is their average number of major mergers, with fast rotators having undergone one major merger, while slow rotators have undergone two. We use this fact and the merging model developed in the previous section, to predict the fraction of slow rotators in the local universe, f_{slow} . We simply assumed that wet major mergers produce fast rotators (first major merger event), dry major mergers produce slow rotators (second major merger event), and dry minor mergers do not change the kinematical state of ETGs. In addition, we took all the ETGs at $z = 1.3$ as fast rotators.

With the previous assumptions, we expect $f_{\text{slow}} \sim 60\%$, in good agreement with the ATLAS^{3D} result of $f_{\text{slow}} \sim 47 - 75\%$ for ETGs with a dynamical mass $M_{\text{dyn}} \geq 10^{11.25} M_{\odot}$ (Emsellem et al. 2011). Thus, our model also reproduces the local fraction of slow/fast rotators of massive ETGs galaxies, reinforcing the conclusions of the previous section.

6.4.2. Uncertainties in the model

The model presented in previous sections has set most of its parameters from observational results. However, there are two unconstrained parameters that could affect our conclusions. The first parameter is f_{disc} , the fraction of gas-rich major mergers whose remnant rebuild a disc component and do not contribute to the increase in ρ_{ETG} . We assumed $f_{\text{disc}} = 0$, and in the following we justify this selection. Hopkins et al. (2009) find that a high gas fraction prevents the destruction of the disc component after a major merger in their N -body/hydrodynamical simulations (but see Bournaud et al. 2011 for a different point of view). These simulations suggest that disc rebuild could be an efficient process when the gas fraction is $f_{\text{gas}} \gtrsim 50\%$. However, we have focused in the massive end of the galaxy population, where the gas fractions are lower. Thanks to Eq. (20), we can estimate the gas fraction of the gas-rich mergers in our model. We find $f_{\text{gas}} \lesssim 30\%$, justifying the assumed $f_{\text{disc}} = 0$. At lower masses than explored in the present section, the gas fraction is higher, and disc rebuild could be an important process in the formation of bulge-dominated spirals (e.g., Huertas-Company et al. 2010; Puech et al. 2012). However, we cannot discard positive values of f_{disc} for massive galaxies, as we see below.

The second parameter is the assumed merger time scale, which typically has a factor of two uncertainty in their normalization (e.g., Hopkins et al. 2010). The T_{MM} from Kitzbichler & White (2008) are typically longer than others in the literature (e.g., Patton & Atfield 2008; Lin et al. 2010) or similar to those from N -body/hydrodynamical simulations (Lotz et al. 2010b,a). Thus, we expect, if anything, a shorter T_{MM} , which implies a larger number density of ETGs due to mergers. Nevertheless, the good description of the ρ_{ETG} evolution with our merger model strongly suggests that mergers are indeed the main process in massive ETGs assembly. Thus, a lower T_{MM} (i.e., a higher merger rate that translates to an ex-

cess of ETGs) could be compensated by a positive value of f_{disc} , that reduces the number density of ETGs due to mergers.

Future observational studies will be important to better constraint the parameters of our model, and further theoretical efforts are needed to understand the uncertainties in the assumed parameters.

7. Summary and conclusions

Using SINFONI/VLT 3D spectroscopy, we have been able to measure, for the first time with spectroscopically-confirmed close pairs, the gas-rich major merger fraction and merger rate at around the peak in star formation activity at $0.9 < z < 1.8$, from the MASSIV sample of star-forming galaxies with a stellar mass range $M_{\star} = 10^9 - 10^{11} M_{\odot}$. In this redshift range we identify 20 close pairs, and classify 13 as major mergers based on a separation $r_p \leq 20h^{-1} - 30h^{-1}$ kpc and a relative velocity $\Delta v \leq 500$ km s⁻¹.

We find that the gas-rich major merger fraction is high, $20.8^{+15.2}_{-6.8}\%$, $20.1^{+8.0}_{-5.1}\%$, and $22.0^{+13.7}_{-7.3}\%$ for $r_p \leq 20h^{-1}$ kpc close pairs in redshift ranges $z = [0.94, 1.06]$, $[1.2, 1.5]$, and $[1.5, 1.8]$, respectively. When compared to measurements at redshifts $z < 1$, the evolution of the (gas-rich) merger fraction can be parametrised as $f_{\text{MM}} = 0.0066 \times (1+z)^m$ with $m = 3.91 \pm 0.16$. We note that the evolution between $z = 1$ and $z \sim 1.5$ seems to flatten out compared to lower redshifts.

The merger rate has been derived using merger time scales from the literature: we find that the gas-rich merger rate is $0.116^{+0.084}_{-0.038}$ Gyr⁻¹, $0.147^{+0.058}_{-0.037}$ Gyr⁻¹, and $0.127^{+0.079}_{-0.042}$ Gyr⁻¹ at $z = 1.03, 1.32$, and 1.54 , respectively, for merger time scales of $T_{\text{MM}} \sim 1.5$ Gyr. We then find that the (gas-rich) merger rate's evolution for galaxies with stellar mass $\bar{M}_{\star} = 10^{10-10.5} M_{\odot}$ over $z = [0, 1.5]$ scales as $(1+z)^n$ with $n = 3.95 \pm 0.12$.

Using these measurements, we developed a simple model to estimate the contribution of gas-rich major mergers to the growth of galaxies on the red sequence. We infer that $\sim 35\%$ of the star-forming galaxies with stellar mass $\bar{M}_{\star} = 10^{10} - 10^{10.5} M_{\odot}$ have undergone a major merger since $z \sim 1.5$. The number of major merger events was about 5 times higher at $1 < z < 1.5$ compared to $z < 1$. Assuming that each gas-rich major merger produces a new early-type galaxy, we infer that the number of gas-rich major mergers is large enough at $z > 1$ to explain the increase in the number density of massive ETGs, supporting a picture where gas-rich (wet) merging is the main process building-up the red sequence. While gas-rich mergers become rarer towards lower redshifts, the number of dry mergers is steadily increasing, and the combination of these two processes accounts for most, if not all, of the increase in the number density of massive ETGs since $z \sim 1.3$. Two-thirds of this number density evolution is due to wet major mergers, while one-third is coming from major and minor dry mergers. These results add further evidence to a picture where merging is a major process driving mass assembly into the massive red sequence galaxies.

We note that minor merging is definitely present in the MASSIV sample (see Sect. 4). However, due to incompleteness in detecting these faint companions, we are not able to assess a minor merger rate at these epochs from our data. In the global picture of red sequence assembly, we emphasise that a simple extrapolation of the minor merger rate measured up to $z \sim 1$ by López-Sanjuan et al. (2011), would indicate that from $z \sim 1.5$ to the present, minor mergers with $1/10 \leq \mu < 1/4$ are not common enough to significantly move spiral galaxies into the red sequence.

⁹ <http://www-astro.physics.ox.ac.uk/atlas3d/>

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To get a complete picture of the life-time effect of major merging on massive galaxies, accurate measurements of the merger fraction and merger rate are needed beyond $z \sim 2$. Spectroscopic surveys will remain an important element to provide secure identification of merging systems at these early epochs.

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Appendix A: Spectroscopic Success Rate in the VVDS-Wide fields

We computed the Spectroscopic Success Rate (*SSR*) in the VVDS-Wide 14h and 22h fields for a given redshift range by following the prescriptions in Ilbert et al. (2006): we compared

the number of galaxies with flag = 4, 3 and 2 over the total number of galaxies (those with flag = 4, 3, 2, 1 and 0). Because flag = 1 galaxies have a 50% reliability and no redshift information is available for flag = 0 galaxies, we took advantage of the latest photometric redshifts from CFHTLS survey to estimate the number of galaxies with flag = 1 and 0 that belong to the redshift range of interest. With the previous steps we estimate the SSR in the 22h field as

$$SSR_{22h}(z_r) = \frac{N_{22h,spec}(z_r)}{N_{22h,spec}(z_r) + N_{22h,phot}(z_r)}, \quad (A.1)$$

where $N_{22h,spec}(z_r)$ is the number of galaxies with flag = 4, 3 and 2 in the redshift range z_r , and $N_{22h,phot}(z_r)$ is the number of galaxies with flag = 1 and 0, and with a photometric redshift in the same redshift range. As the available imaging data in the 14h field is not as deep as the CFHTLS, we show in the following that we can assume that the SSR in the 14h field follows the same distribution as in the 22h field, for which CFHTLS photometry is available.

First we checked the properties of the stars in these two VVDS-Wide fields. The fraction of stars in the 22h (14h) field is 66% (51%) for flag = 4 sources, 38% (26%) for flag = 3 sources, 24% (23%) for flag = 2 sources, and 11% (12%) for flag = 1 sources. So the fraction of stars is similar in both fields for flag = 2 and 1 sources, while higher in the 22h field for flag = 4 and 3. We checked that the normalised distributions of the VVDS-Wide stars as a function of their observed I_{AB} -band magnitude in both fields are similar for each flag. We note that stars with high confidence flag are brighter than the low confidence ones, as expected. The fraction of bright stars ($I_{AB} \leq 21$) in the 22h (14h) field is 75% (67%) for flag = 4 stars, 49% (43%) for flag = 3 stars, 30% (29%) for flag = 2 stars, and 19% (16%) for flag = 1 stars. These fractions suggest that there is a higher density of bright stars in the 22h area than in the 14h one, that translates in a higher fraction of stars for flag = 4 and 3 sources, while faint stars have similar densities in both fields, leading to a similar fraction of stars with flag = 2 and 1. Because of this, we assumed that the fraction of stars among flag = 0 sources, that we estimate in the 22h field from the CFHTLS photometry, is similar in 22h and 14h fields.

The distribution of galaxies at $z \geq 0.9$, the redshift range in which we are interested on, are also similar in both fields. Because of this, we assumed that the photometric distribution of galaxies with flag = 1 and 0 is similar in both fields, and we use those in 22h field to estimate $N_{14h,phot}$ and the SSR in the 14h field:

$$SSR_{14h}(z_r) = \frac{N_{14h,spec}(z_r)}{N_{14h,spec}(z_r) + f_{22h}(z_r) \times n_{14h}(1, 0)}, \quad (A.2)$$

where $n_{14h}(1, 0)$ is the total number of sources (galaxies and stars) with flag = 1 and 0 in the 14h field, and $f_{22h}(z_r) = N_{22h,phot}(z_r)/n_{22h}(1, 0)$ is the fraction of sources with a photometric redshift in the redshift range z_r over the total population of sources with flag = 1 and 0 in the 22h field.

Appendix B: GALFIT residual images of blended MASSIV close pairs

In this Appendix we detail the modelling in the i band of those MASSIV close pair candidates with overlapping components in the kinematical maps. We used GALFIT v3.0 (Peng et al. 2010) to model the light distribution with two independent Sérsic components. We set the initial positions of the sources using the information from the kinematical maps, and we did not impose any

constraint on the other initial parameters of the fit (i.e., luminosity, effective radius, Sérsic index, position angle and inclination). Because of the minimisation process preformed by GALFIT in the fitting, the best model with two components should not be unique and the convergence to a given solution should depend on the initial values of the parameters defined by the user. We checked, by exploring randomly the space of initial values, that the number of good solutions is at most two. Where two good solutions exist, the one that better reproduces both the velocity field and the velocity dispersion map is preferred. To illustrate the need of a second component to describe the blended close pair systems in MASSIV, we show in Fig. B.1 the original image in the i band and the residual image from GALFIT with one and two components. In all the cases the residual map from the one component fit suggests that a second component is present. This Figure also demonstrates that both the seeing ($\sim 0.7''$) and the depth of the i -band images used in the present study are good enough to characterise sources with two overlapping components.

Appendix C: Completeness of the close pair sample

Throughout the present paper we have assumed that the close pair systems detected in the MASSIV sample are representative for galaxies with $I_{AB} \leq 23.9$. The MASSIV sample comprises galaxies fainter than this luminosity limit, and in this appendix we justify the boundary applied in our analysis.

The VVDS parent samples of the MASSIV sample are randomly selected in the I_{AB} band (Le Fèvre et al. 2005). Because of this, we studied the distribution of MASSIV sources as a function of the I_{AB} -band magnitude to obtain clues about the completeness of the MASSIV sample in our close pair study. We show the cumulative distribution of all MASSIV sources and of those with a close companion (both major and no major) in the top panel of Fig. C.1. Obviously, we are able to detect single MASSIV galaxies at fainter magnitudes than the close pairs ($I_{AB} = 24.4$ vs 23.8) because in the latter case we have to detect both the principal source and the companion galaxy, which is usually fainter.

Lets assume for a moment that the detection curve of the MASSIV sources is a step function that is 1 for $I_{AB} \leq I_{AB,lim}$ and 0 for fainter magnitudes. In this ideal case, $I_{AB,lim}$ is defined by the fainter galaxy (close pair) detected. Thus, we should only be able to detect close pairs at $I_{AB} \leq 23.8$, and the measured merger fraction in the total MASSIV sample ($I_{AB} \leq 24.4$) will be lower than the real merger fraction because of the missing faint close pairs with $I_{AB} > 23.8$. Of course, the detection curve in the MASSIV sample is more complicated than a step function, and could depend on redshift, geometry, luminosity, etc. Instead of trying to estimate the detection curve for MASSIV sources and for close companions to recover statistically the missing close pairs (see Patton et al. 2000; de Ravel et al. 2009, for examples of this kind of correction), we define the I_{AB} -band magnitude up to which the detected close pairs are representative of the total MASSIV sample, named $I_{AB,comp}$, and in our study we only keep those sources brighter than $I_{AB,comp}$ to ensure reliable merger fractions.

The MASSIV sample is a representative subsample of the global star-forming population (Contini et al. 2012), and we expect the MASSIV galaxy pairs to be also a random sample of this global population. Therefore, the distributions in the I_{AB} band of the total and the close pair MASSIV sources should be similar when the close pairs detection was slightly affected by incompleteness issues. Following this idea, we performed

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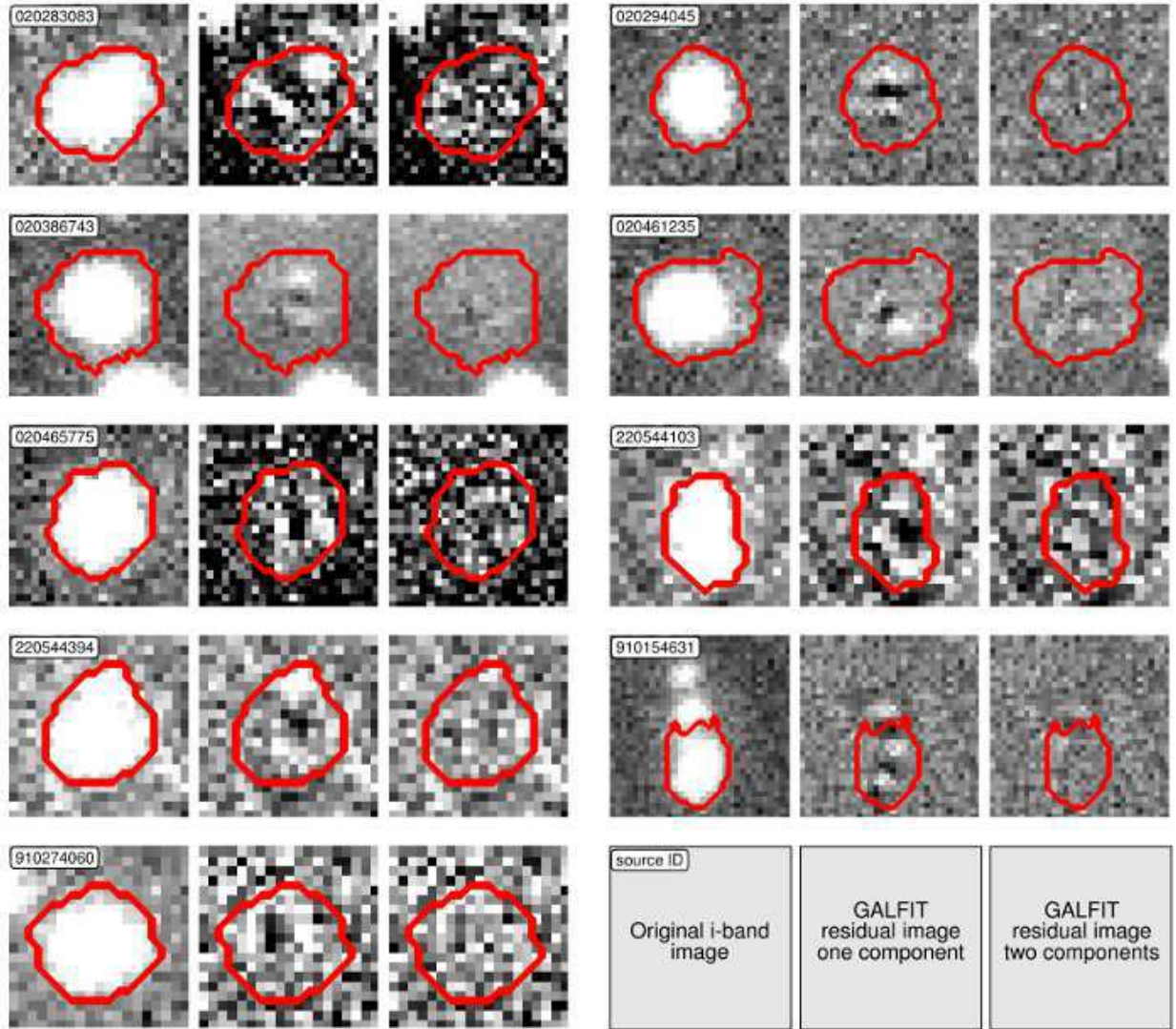


Fig. B.1. Original i -band image (*left panels*) and residual images from the GALFIT fit with one (*central panels*) and two components (*right panels*) of those MASSIV close pair candidates with overlapping components in the kinematical maps. The grey scale has been chosen to enhance the light residuals in the images. The solid contour marks the targeted MASSIV source to guide the eye. The ID of the showed MASSIV source is labelled in each left panel. [A colour version of this plot is available at the electronic edition].

a Kolmogorov-Smirnov (KS) test over the total and close pair MASSIV galaxies with $I_{AB} \leq I_{AB,lim}$, and explored different values of $I_{AB,lim}$, from 23.0 to 24.5. Then, the completeness magnitude $I_{AB,comp}$ was defined by the minimum in the KS estimator D , i.e., where both distributions have the lower probability to be different. This exercise provides the curve in the bottom panel of Fig. C.1, that states $I_{AB,comp} = 23.9$. We repeated this procedure with the major merger sample, and we obtain a similar distribution of D values, reinforcing our choice.

In summary, in the estimation of the merger fraction we only use those MASSIV galaxies with $I_{AB} \leq I_{AB,comp} = 23.9$. This ensures that the close pair sample is a random subsample of the total MASSIV one, and we avoid any bias related with the shal-

lower detection curve of close pairs compared with that of the total MASSIV sample.

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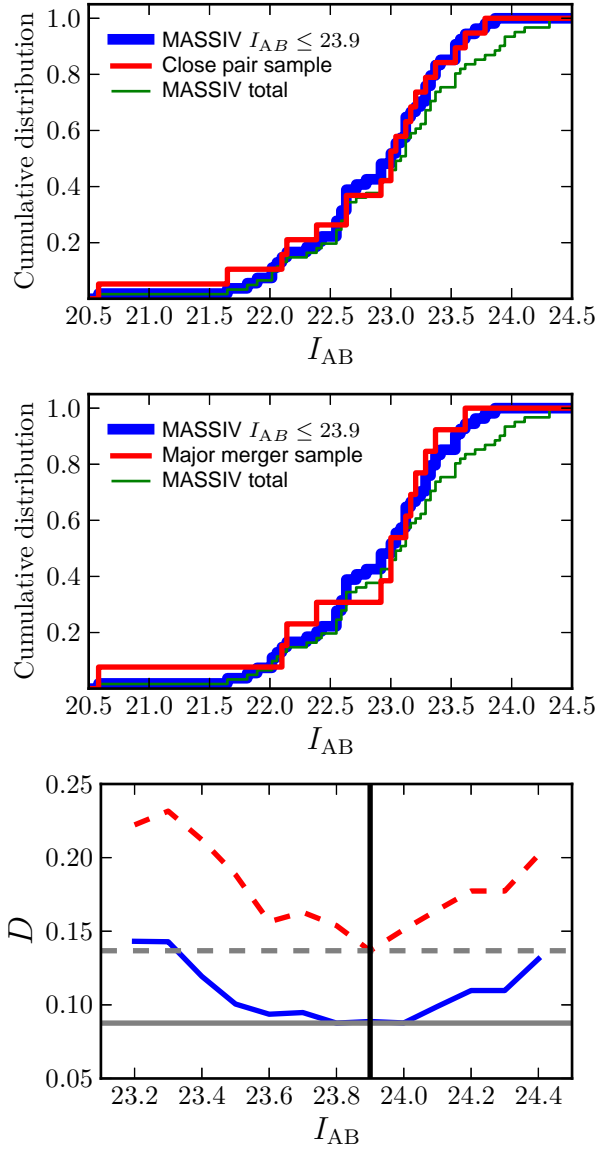


Fig. C.1. *Top:* cumulative distribution in the apparent I_{AB} magnitude of the total MASSIV sample (green line), the close pair sample (red line) and the total MASSIV sample with $I_{AB} \leq 23.9$ (blue line). *Middle:* the same than before, but the red line shows the distribution of the major merger sample. *Bottom:* Kolmogorov-Smirnov estimator D of the total MASSIV sample distribution against the close pair (blue solid line) and the major merger (red dashed line) distributions. We computed D for those galaxies brighter than a given I_{AB} magnitude. The solid (dashed) horizontal line marks the minimum in the D distribution for close pairs (major mergers). The vertical solid line marks the I_{AB} magnitude at the minimum value of D , $I_{AB,comp} = 23.9$. [A colour version of this plot is available at the electronic edition].

Programmes

Résumé

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Cette annexe présente les trois principaux programmes informatiques développés au cours de ce travail de thèse. Pour chacun d'entre eux, je détaille les pré-requis d'installation ainsi que les paramètres d'entrées nécessaires à leur fonctionnement. Ces trois programmes ont vocation à être diffusés dans la communauté scientifique, et sont donc tous les trois libre d'accès sur internet.

C.1 DICE

C.1.1 Purpose

DICE is designed to generate initial conditions file, allowing the user to explore the dynamics (and many other physics) of galaxies through secular evolution, or through encounter events. Generating idealized galactic initial conditions involves using a lot of different numerical tools and algorithms. Thus, reading and understanding the code should never be considered as a waste of time. DICE intends to be as much didactic as starscream was for me. With this code, you can generate initial conditions using only ascii files as input. The advantage of DICE is that you can generate with one executable a lot of different scenarios.

C.1.2 Required libraries

DICE uses the GSL & FFTW-3 libraries. You can find these libraries here :

- <http://www.fftw.org/download.html>
- <http://www.gnu.org/software/gsl/>

GSL is used to generate random numbers, and also to perform numerical integration. FFTW is used to compute Fourier Transforms through the fastest portable implementation available.

On Linux-based systems, it is quite easy to perform a standard installation of these libraries using package tools such as **apt-get** or **yum**. If you are working on a HPC cluster, you may be able to load these libraries using the `module load library` command.

On OSX systems, **macport** (<http://www.macports.org>) is a good option to compile and install easily these libraries. You might be interested in using the threading capacities of the library FFTW. In this case, download the tarball, and perform the following commands :

```
tar -xzvf fftw-x.x.x.tar.gz
cd fftw-x.x.x
./configure --enable-threads
make
sudo make install
```

If your libraries are installed in a non-standard location, think about updating the `DYLD_LIBRARY_PATH` variable for shared libraries and `LD_LIBRARY_PATH` variable for static libraries.

C.1.3 Compilation & installation

DICE is available for download from the internet. You can checkout the latest version in the DICE SVN repository typing :

```
mkdir DICE
cd DICE
svn checkout http://dice-project.googlecode.com/svn/trunk/
```

The DICE package comes with the CMake cross-platform build system. So technically, you don't have to worry so much about the compilation. Make sure you have cmake installed by typing :

```
cmake --version
```

If your version is older than cmake 2.6, you will have to update your system to a more recent version of cmake. To generate the makefile, type :

```
cd build
cmake ..
```

If no errors are detected through this step, compile the code and install the code like this :

```
make
make install
```

By default, DICE is installed in `$HOME/local/dice`, but you can specify a different installation directory using the flag `-DCMAKE_INSTALL_PREFIX=/install/path`

A `cmake` macro is implemented to locate standard installations of these libraries. Nevertheless, if you installed them in a different way, use the following keywords to help `cmake` locating these libraries :

- `-DGSL_PATH = /path/to/gsl`
- `-DFFTW3_PATH = /path/to/fftw3`

Once installed, think about adding to the `PATH` environment variable the location of the DICE binary. If you use `bash` and under standard installation process, add to your `.bashrc` (or `.profile` for OSX systems) :

```
export PATH=$HOME/local/dice/bin:$PATH
```

Optionally, if you want the code to run a bit faster, you can use the threading abilities of the FFTW library for the gravitational potential computation. In this case, and assuming that you have installed the treaded version of FFTW3, then replace the previous commands by :

```
cmake .. -DENABLE-THREADS=ON
make
make install
```

CMake will automatically look for the `fftw3_threads` library, and link it.

C.1.4 Quick start

DICE take as input the name of a file called the DICE configuration file. This file contains all the informations needed to start the program. It specifies :

- The location of ASCII files containing the physical parameters of the galaxy to generate. These files are called the galaxy parameters files. You can point up to 64 galaxy parameters files , just remember to put galaxies with same parameters one after the other : it lowers the computation time since DICE copy the result of the previous computation if it is strictly the same galaxy.
- The output format of the initial conditions. For now, DICE handle only two format. **Gadget1** and **RamsesGroup**. But you can use the `uns_projects` tools, calling the `unsio` API, to convert from **Gadget1** to **nemo** or **Gadget2**.
- The parameters to define a Keplerian trajectory between two and only two galaxies. If there more or less galaxies, or if DICE misses one or more of the three keywords needed, the galaxies will not be placed on such an orbit.

DICE is provided with an example, in order to quickly test the DICE initial conditions. Once DICE has been installed, you can run the example by typing in the prompt :

```
cd example
dice dice.config
```

Contents of "dice.config" file in example folder :

```
% List of galaxy parameters file
% Putting identical galaxies one beside the other allows DICE
% to re-use the previous computation,
% and thus reduces the global execution time
Galaxy ./galaxy1.params
Galaxy ./galaxy2.params

% Output format of the initial conditions
% Allowed format 'Gadget1','RamsesGroup'
ICformat Gadget1

% Keplerian trajectory parameters
% This keywords should be set
% if you want a Keplerian orbit between two and only two galaxies!!
% If there are more or less galaxies, trajectories will be set according
% to the xc,yc,zc,vel_xc,vel_yc,vel_zc keywords in the galaxy parameters files.
% It induces that the keywords xc,yc,zc,vel_xc,vel_yc,vel_zc are not considered
% when the Keplerian trajectory parameters are set.
% By default, galaxies are aligned along the X-axis

% Eccentricity of the Keplerian orbits
Eccentricity 1.0
% Initial distance between the galaxies [kpc]
Rinit 35.
% Pericentral distance of the Keplerian system [kpc]
Rperi 10.
% Polar angle of the normal vector of the orbital plane
OrbitPlanePhi 45.
% Azimuthal angle pf the normal vector of the orbital plane
OrbitPlaneTheta 45.

Nthreads 4
GasHydrostaticEq 1
MeanPartDist 0
AcceptImaginary 1
```

Each Galaxy keywords in the dice configuration file should point to an ASCII file containing the physical parameters of the galaxy to construct. Every keywords are needed, so check the

content of this file carefully.

Content of galaxy1.params file in example folder :

```
% Virial velocity of the galaxy [km/s]
v200 200.0
% Redshit of the galaxy
redshift 0.00
% Mass fraction of the baryons disk
m_d 0.05
% Disk circular velocity fraction
f_d 1.00
% Bulge circular velocity fraction
f_b 0.10
% Disk angular momentum fraction
j_d 0.05
% Bulge-Disk mass ratio
bd_ratio 0.10
% Halo spin parameter
lambda 0.04
% Halo concentration parameter [kpc]
c_halo 9.00
% Stellar disk scale length [kpc]
% If the value is negative, the scalelength will be computed
% using the fitting formula provided in Mo et al. 1998
r_disk 4.00
% Gaseous disk scale length [kpc]
r_gas 4.00
% Bulge scale length [kpc]
r_bulge 0.50
% Number of element for the potential grid
% Setting ngrid as 2^n with n positive integer
% allows FFTW algorithm to run much more faster
level_grid 8
% Size of the box to compute the gravitational potential [kpc]
boxsize 160.0
% Mass fraction of gas compare to stellar disk
gas_fraction 0.15
% Initial temperature of the gas [K]
t_init 1e4
% Stellar disk thickness parameter
disk_flat 0.15
% Gaseous disk thickness parameter
gas_flat 0.05
...
```

```
...
% Halo flattening
halo_flat 1.00
% Bulge Flattening
bulge_flat 0.70
% Minimal value for the Toomre parameter
Q_lim 0.0
% Number of halo particles
nhalo 100000
% Number of stellar disk particles
ndisk 100000
% Number of gaseous disk particles
ngas 100000
% Number of stellar bulge particles
nbulge 0
% Halo radial density cut [kpc]
halo_cut 24.5
% Stellar disk radial density cut [kpc]
disk_cut 12.0
% Gaseous disk radial density cut [kpc]
gas_cut 12.0
% Stellar bulge density cut [kpc]
bulge_cut 1.0
% Axisymmetric_drift approximation in the Jeans equations
axisymmetric_drift 0
% Dispersion extinction coefficient. Values need to be in the interval [0,1].
% For example, setting DispExtCoeff to 0.90
% means that the azimuthal dispersion is equal to 90 percent
% the value found by integrating the Jeans equations
% at a radius equal to the disk scale length.
% The function used to extinguished is (1-exp(-r/rh))
% If this coefficient is outside the range [0,1],
% no extinction is applied to the radial an azimuthal velocity dispersion
DispExtCoeff 0.90
% Maximum velocity of a disk particle in expressed
% in units of escape velocity
disk_vmax 3.5
% Maximum velocity of a halo particle in expressed
% in units of escape velocity
halo_vmax 3.5
% Maximum velocity of a bulge particle in expressed
% in units of escape velocity
bulge_vmax 3.5
% Gaussian steps for the MCMC-MH particle positioning algorithm,
% expressed in units of the component scale length. Default value is 0.5.
disk_mcmc_step 0.50
gas_mcmc_step 0.50
halo_mcmc_step 0.20
bulge_mcmc_step 0.50
...
```

```
...
% Halo model
% 1 = Hernquist profile
% 2 = Plummer profile
% 3 = Jaffe profile
% 4 = Uniform profile
% 5 = NFW profile
halo_model 1
% Disk model
% 1 = Exponential-r Sech2-z disk
% 2 = Myamoto-Nagai profile
% 3 = Exponential-r Exponential-z disk
% 4 = Hernquist profile
disk_model 2
% Gas model
% 1 = Exponential-r Sech2-z disk
% 2 = Myamoto-Nagai profile
% 3 = Exponential-r Exponential-z disk
gas_model 3
% Bulge model
% 1 = Hernquist profile
% 2 = Plummer profile,
% 3 = Jaffe profile
bulge_model 1
% Position of the galactic center on x-axis [kpc]
xc 0.
% Position of the galactic center on y-axis [kpc]
yc 0.
% Position of the galactic center on z-axis [kpc]
zc 0.
% X-axis velocity component of the galaxy [km/s]
vel_xc 0.
% Y-axis velocity component of the galaxy [km/s]
vel_yc 0.
% Z-axis velocity component of the galaxy [km/s]
vel_zc 0.
% Spin angle of the disk in the XY plane [degree]
spin 0.
% Inclination of the disk compare to XY plane [degree]
incl 0.
% Seed for the random number generator
seed 123456
```

C.2 AURORA

Aurora is a python program which allows to create :

1. SPH (or AMR using pseudo-SPH approximation) and N-body projected maps,
2. photometry mock observations,
3. 3D spectroscopic mock observations,

using numerical simulations of galaxies stored in the Gadget2 file format.

```
mkdir aurora
cd aurora
svn checkout http://aurora-py.googlecode.com/svn/trunk/
```

To run Aurora, you need to add the location of the `aurora.py` file to your `PYTHONPATH` environment variable. Then, type into the prompt :

```
python
import aurora
aurora.maps('aurora.conf')
aurora.photom_mock('aurora.conf')
aurora.spectrom_mock('aurora.conf')
```

The Aurora configuration file `aurora.conf` file must contain the following keywords to be able to create the desired maps :

```
# Parameters for the creation of a spectroscopic 3D mock cube.
[spectrom]
# Presets for all the 'spectrom' parameters.
# Available presets: 'sinfoni', 'sinfoni-ao', 'kmos', 'eagle', 'muse-wide',
# 'muse-narrow', 'ghasp-instr'
presets = sinfoni
# Spatial sampling of the mock cube [arcsec].
spatial_sampl = 0.125
# Spectral sampling of the mock cube [Angstrom].
spectral_sampl = 1.95
# Spatial resolution of the mock cube [arcsec].
spatial_res = 0.65
# Spectral resolution of the mock cube.
spectral_res = 2500
...
```

```
...
# Number of spatial elements of the mock cube.
spatial_dim = 38
# Number of spectral elements of the mock cube.
spectral_dim = 48
# Mean dispersion of the noise [10e-16 erg.s^-1.cm^-2.micron^-1].
# This option defines the noise level.
sigma_cont = 5.0
# Targeted best SNR in the mock cube.
# This alternative option defines the noise level.
# If defined, it overrides the sigma_cont noise definition
target_snr = 0

# Parameters for the creation of hydrodynamical quantities maps.
[maps]
# Presets for all the 'maps' parameters.
presets = sdss
# Number of spatial elements of the maps.
spatial_dim = 512
# Spatial sampling of the mock cube [arcsec].
spatial_sampl = 0.396
# Method used to compute the hydrodynamical quantities.
# 'mean' enables a mass weighed computation,
# while 'max' computes the maximum value along the line-of-sight
method = mean

# Parameters for the creation of spectrometric mock maps.
[photom]
# Presets for all the 'maps' parameters.
# Available presets: sdss_r, sdss_u, sdss_g, sdss_r, sdss_i, sdss_z,
# cfht_mega_i, hst_wfc3_f160w, hst_acs_f814w
presets = sdss_r
# File containing the SED model.
# It must formatted as a standard Starburst99 file:
# Time [yr] Wavelength [Ang] &
# Log(total lum) Log(stellar lum) Log(nebular lum) [erg/sec/a]
sed_file = salpeter_S99.spectrum1
# Reference mass of the stellar population modeled by the SED
sed_reference_mass = 1E4
# Transmission of filters.
# This parameter can be a list of files.
# Aurora will loop over this list, and produce a mock observation for each of them
filter_file = SDSS_U.dat,SDSS_G.dat,SDSS_R.dat
...
```

```
...
# Size of the stellar particles packets.
# Values in the format 2^n with n a positive integer should be preferred.
# Aurora sort the stellar particles by their ages,
# and split the list of particles into packet of size 'nvector_stars'.
# Then, Aurora assign the same SED in a given packet.
nvector_stars = 32
# Magnitude limit expressed in AB system.
# This value allows to compute the noise level.
mag_ab_limit = 25.

# Orientation parameters of the simulation
[geometry]
# Spectral redshift.
# From this value, Aurora computes a cosmological luminous distance
# according to Planck-2013 cosmology.
redshift = 0.01435
# Polar angle for the rotation matrix [degree]
theta = 54.3301
# Azimuthal angle for the rotation matrix [degree]
phi = -43.3253
# Optional reentering of the simulation to its barycenter
barycenter = false
# Scale factor for the spatial quantities of the simulation.
scale = 1.0
# Final size of the pixel in any mock/map output [kpc].
# This parameters overrides the 'redshift' parameter,
# and skip the cosmological luminous distance computation. Ignored if set to 0.
pixsize = 0.0
# Manual center of the Cartesian x component of the simulation [kpc]
centerx = 0.150000
# Manual center of the Cartesian y component of the simulation [kpc]
centery = -0.250000
# Manual center of the Cartesian z component of the simulation [kpc]
centerz = 0.150000
# Line-of-sight selection.
# Before applying the rotation matrix to the simulation,
# this parameter select the Cartesian axis,
# which should be used as the line of sight.
# Available: x,y,z.
los = z
...
```

```
...
# Aurora run parameters
[run]
# Simulation file. The file must be formatted with the GADGET2 format.
input_file = output_00200.g2
# Minimum density to consider for the mock/map computation [cm-3].
rho_min = -1E-5
# Maximum density to consider for the mock/map computation [cm-3].
rho_max = 1E10
# Size of the stellar particles packets. Higher
# Aurora works with packets of gas Lagrangian particles,
# usin order to avoid sequential code, which is much slower in python.
# Values in the format 2n with n a positive integer should be preferred.
# High values should make the run faster,
# but requires larger amount of memory
nvector = 2097152
# Number of kernels for the Gaussian smoothing of the gas particles.
# Aurora sort the gas particles in families of smoothing length,
# and smooth each family with a kernel proportional
# to the mean smoothing length of the family.
nfft = 9
# Size of the smallest Gaussien kernel used for the family of gas particles
# with smallest smoothing length [pc].
# This parameter determines the kernel size for all the other 'nfft' families.
fft_hsml_min = 7.5

# Properties of the simulation
[input_properties]
# Units system of simulation output
# Available: ramses, gadget2
units = ramses
```


C.3 KINEGEN

KINEGEN (KINemetry in GENetic algorithm) is an IDL routine that enables to adjust the geometrical parameters of a velocity field using a genetic algorithm strategy. For each set of geometrical parameters, KINEGEN models a velocity field using the rotation harmonic coefficient of the kinemetry method. You can checkout the latest version in the KINEGEN SVN repository typing :

```
mkdir kinegen
cd kinegen
svn checkout http://kinegen.googlecode.com/svn/trunk/
```

KINEGEN is an IDL procedure. The parameters of this procedure are the following :

The minimal command line in the IDL prompt to get KINEGEN adjust geometrical parameters on a velocity field contained in a FITS image "rv.fits" is :

```
kinegen, 'rv.fits'
```

The full KINEGEN parameters are listed hereafter :

```
; Mandatory Input
;
; FITS file that contains the velocity field to be fitted
file = 'velocity_field.fits'

; Optional inputs
;
; Measurement of errors on the radial velocity.
error_file = error_file
; Signal-to-noise map corresponding to the input velocity field
; This keyword overrides the previous 'error_file' keyword.
snr_file = snr_file
...
```

```
...
; PARINFO: An IDL structure, that defines
; the parameters geometrical parameters limits.
; This structure is strictly similar to the one used by the MPFIT IDL routine.
; See the MPFIT documentation to see the entries of the par info structure.
parinfo = parinfo

; PLOTALL: Plot to the screen all the generated kinemetry velocity fields
plotall = plotall

; PLOTLAST: Plot to the screen the final kinemetry model
plotlast = plotlast

; VERBOSE: Verbose mode
verbose = verbose

; NAME: Prefix name for the output file
; that stores the geometrical parameters values
name = 'rv'

; SEED: Seed for the random number generator
; If not specified, the seed is initialized
; with the internal clock function 'systemtime'
seed = 11

; CTRL: Genetic algorithm parameters
; Array with 13 values:
;   ctrl[ 1] - number of individuals in a population (defaultabst
;             is 100)
;   ctrl[ 2] - number of generations over which solution is
;             to evolve (default is 500)
;   ctrl[ 3] - number of significant digits (i.e., number of
;             genes) retained in chromosomal encoding (default
;             is 6) (Note: This number is limited by the
;             machine floating point precision. Most 32-bit
;             floating point representations have only 6 full
;             digits of precision. To achieve greater preci-
;             sion this routine could be converted to double
;             precision, but note that this would also require
;             a double precision random number generator, which
;             likely would not have more than 9 digits of
;             precision if it used 4-byte integers internally.)
;   ctrl[ 4] - crossover probability; must be <= 1.0 (default
;             is 0.85)
;   ctrl[ 5] - mutation mode; 1/2=steady/variable (default is 2)
...
```

```

...
; ctrl[ 6] - initial mutation rate; should be small (default
;           is 0.005) (Note: the mutation rate is the proba-
;           bility that any one gene locus will mutate in
;           any one generation.)
; ctrl[ 7] - minimum mutation rate; must be >= 0.0 (default
;           is 0.0005)
; ctrl[ 8] - maximum mutation rate; must be <= 1.0 (default
;           is 0.25)
; ctrl[ 9] - relative fitness differential; range from 0
;           (none) to 1 (maximum). (default is 1.)
; ctrl[10] - reproduction plan; 1/2/3=Full generational
;           replacement/Steady-state-replace-random/Steady-
;           state-replace-worst (default is 3)
; ctrl[11] - elitism flag; 0/1=off/on (default is 0)
;           (Applies only to reproduction plans 1 and 2)
; ctrl[12] - printed output 0/1/2/3=None/Minimal/
;           Verbose/Histogram
;           (default is 0)
ctrl = [300,30,8,.85,2,.20,.20,.75,1.0,3,1,1]

; THREADING: Launch multiple threads through
; the IDLBridge interface
; Still experimental. The number of threads is determined
; thanks to the environment variable !CPU.TPOOL_NTHREADS
/threading

; SEEING: Point Spread Function full width at half maximum expressed in pixels.
seeing = 5.0

; RWD_INDEX1: Fitness boost factor.
; This parameter boost geometrical parameters
; that enables to better cover the velocity field
rwd_index1 = 6.0

; RWD_INDEX2: Fitness boost factor.
; This parameter boost geometrical parameters
; that enables a kinematical center close to the flux center.
; Requires the use of SNR_FILE or ERROR_FILE
; to locate the flux center
rwd_index2 = 6.0

; D0: Estimation of the disk scalelength expressed in pixels.
; Combined with the seeing keyword, it enables a correction
; of the PSF effect on the kinemetry ellipses
d0 = 3.0
...

```

```

...
; COVER: Kinemetry ellipse cover parameter.
;The kinemetry algorithm stops the expansion
; once an ellipses cover less than COVER% the surface if the velocity field.
cover = 0.75

; RC_MAX_VDEP: Maximum velocity for the deprojected B1
; kinemetry coefficient [km.s^-1]
; This parameter prevents solutions with too low inclinations
rc_max_vdep = 1000

; RC_MIN_SLOPE: Minimum slope for the linear regression
; of the radial B1 values [km.s^-1.px^-1]
; This parameters prevents solutions with too flat rotation curves
rc_min_slope = 0.0

; RC_MAX_DELTA: Maximum step between two
; successive B1 coefficients [km.s^-1].
; This parameter prevents geometrical parameters
; with strong oscillations in the final rotation curve
rc_max_delta = 100.

; RC_MAX_VR0: Maximum value for the B1 term
; adjusted on the first ellipse [km.s^-1].
rc_max_vr0 = 200.

; Output
;
; PARAMS: An array with dimension [4,3] that specifies
; the adjusted geometrical parameters [*,0]
; the upper limit 1 sigma errors [*,1]
; and the lower limit 1 sigma errors [*,2]
params = params

```

To fit the geometrical parameters of a velocity field `rv.fits` with an velocity error map `err_rv.fits`, plotting all the generated velocity field models, with a PSF size equal to 5 pixels on the velocity field, the following command should be invoked in the IDL interpreter:

```
kinegen, 'vel.fits', error_file='err_rv.fits', ctrl=[300,30,8, .85,2, .20, .20, .75,1.0,3,1,1],
```

The custom CTRL input variable allows to generate maintain a population of 300 models per genetic loop, with a crossover rate of 85% and a mutation rate of 20%. At the end of the fit, the geometrical parameters are stored in an ASCII file named `_kinegen_params.dat`. A FITS structure containing the kinemetry coefficients of the harmonic expansion performed with the final set of geometrical parameters is created with the name `_kinegen_coeff.fits`. A prefix can be added to the name of the output file through the optional keyword NAME.

Résumé

La compréhension des mécanismes de formation et d'évolution des galaxies dans un cadre cosmologique demeure un grand défi de l'astrophysique moderne. En particulier, la contribution des processus de fusion lors de l'assemblage cosmologique de masses est encore mal comprise. Dans ce contexte, l'étude du support dynamique des galaxies est un moyen permettant de contraindre les différents scénarios d'évolution. Dans une première partie, je présente ma contribution à l'analyse cinématique des galaxies de l'échantillon MASSIV composé de 83 galaxies jeunes ($1 < z < 2$) observées avec le spectrographe intégral de champ SINFONI au VLT. Grâce à l'utilisation d'une méthode basée sur une décomposition en harmoniques sphériques des champs de vitesses, je définis une technique d'analyse et de classification automatique qui vise à quantifier la fraction d'objets susceptibles d'être le produit d'une fusion récente. J'introduis dans une seconde partie une méthode de construction de conditions initiales de galaxies idéalisées, pré-requis indispensable à la modélisation hydrodynamique de l'interaction de galaxies jeunes. Dans une troisième partie, je présente la définition et l'analyse de l'échantillon de simulations idéalisées MIRAGE, dont les conditions initiales sont conçues de manière à reproduire les propriétés physiques des galaxies jeunes les plus riches en gaz. Cet échantillon qui a mobilisé environ 3,5 millions d'heures CPU, est composé de 20 simulations de fusions explorant les paramètres initiaux de masse et d'orientation des disques avec une résolution spatiale atteignant 7 parsecs (~ 23 années-lumière) et une modélisation de nombreux processus physiques. Les simulations montrent en particulier l'absence de flambées de formation stellaire dans les fusions de disques fragmentés et turbulents, suggérant un mécanisme de saturation. Une analyse des propriétés des complexes granulaires issus d'instabilités gravitationnelles est réalisée dans la quatrième partie de cette thèse, soulignant la compatibilité du scénario granules âgés avec les observations en terme de taux d'éjection de gaz et d'âge de populations stellaires. Je mets en évidence que le mécanisme de fusion de granules riches en gaz peut permettre de contrôler la croissance de la masse du bulbe, d'aplatir le profil central de densité du halo de matière sombre ainsi que d'éjecter de grandes quantités de gaz dans le plan du disque. Enfin, dans une cinquième et dernière partie, je complète ce travail en réalisant une étude comparative des données cinématiques MASSIV à un jeu de plus de 4000 pseudo-observations à $z=1.7$ construites à partir des simulations de l'échantillon MIRAGE afin de déterminer la capacité à détecter des signatures de fusion de galaxies dans le cadre des conditions observationnelles de l'instrument SINFONI.

Mots-clés : Astronomie – Astrophysique – Galaxies – Cosmologie – Simulations numériques – Cinématique – Dynamique – Spectroscopie

Abstract

Understanding the mechanisms of formation and evolution of galaxies in a cosmological context remains a major challenge in modern astrophysics. Especially, the contribution of the fusion processes to the cosmological mass assembly is still poorly understood. In this context, the study of the dynamical support of galaxies is a way to constrain different evolution scenarios. In the first part, I introduce my contribution to the kinematical analysis of galaxies MASSIV sample of 83 young galaxies ($1 < z < 2$) observed with the integral field spectrograph SINFONI at the VLT. Thanks to the use of a method based on a spherical harmonics decomposition of velocity fields, I define an analysis and an automatic classification technique that aims at quantifying the fraction of objects that are likely to be the product of a recent galaxy merger. I introduce in the second part a method to build initial conditions of idealized galaxies, prerequisite to the hydrodynamic modeling of young galaxies in interaction. In the third part, I present the definition and analysis of the sample of idealized simulations MIRAGE, whose initial conditions are designed to reproduce the physical properties of the most gas-rich young galaxies. This sample which required about 3.5 million hours CPU time, is composed of 20 simulations of mergers exploring the initial parameters of mass and orientation of the disks with a spatial resolution reaching 7 parsecs (~ 23 light years) and modeling many physical processes. The simulations show in particular the absence of a star formation bursts in mergers of fragmented and turbulent disks, suggesting a saturation mechanism. An analysis of the properties of clumps that are generated by gravitational instabilities is performed in the fourth part of this thesis, highlighting the compatibility of long-lived clumps scenarios with observations in terms of outflowing material and stellar population ages. I also highlight that the gas rich clump fusion mechanism is able to control the bulge mass growth, to erode the central profile of the dark matter halo and to drive massive gas outflows in the disk plane. Finally, in the fifth and final part, I complete this work by carrying out a comparative study of the MASSIV kinematical data to a set of more than 4000 pseudo-observations at $z=1.7$ built from simulations of the MIRAGE sample to determine the ability to detect galaxy merger signatures under the observational conditions of the SINFONI instrument.

Keywords : Astronomy – Astrophysics – Galaxies – Cosmology – Numerical Simulations – Kinematics – Dynamics – Spectroscopy

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