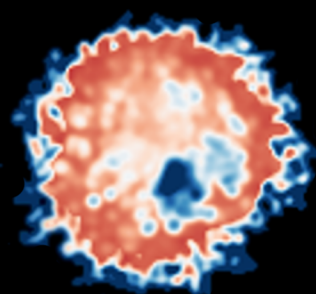
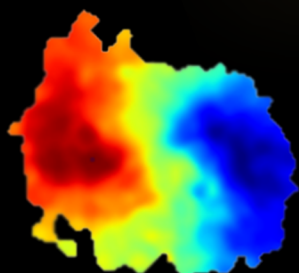


The smallest galaxies: tests for cosmological & galaxy evolution models

De kleinste sterrenstelsels: tests voor kosmologische &
galaxie-evolutie modellen



Robbert Verbeke

Promotor: Prof. Dr. Sven De Rijcke

Proefschrift tot het behalen van de academische graad van
Doctor in de Wetenschappen: Sterrenkunde
Academiejaar 2016-2017

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Dit werk kwam tot stand in het kader van het Interuniversitaire Attractie Polen Programma, opgericht door het Federaal Wetenschapsbeleid (Belspo: IAP P7/08 CHARM)

Figuur voorpagina:

*Boven: de gas distributie van twee interagerende gesimuleerde dwergsterrenstelsels.
Links onder: het snelheidsveld van het gas van een gesimuleerde dwerg (rood duidt
gas aan dat naar ons toe beweegt, in het blauw beweegt het gas zich weg van ons).
Rechts: kleurmap van een gesimuleerde dwerg die een periode van verhoogde ster-
vorming ondergaat. Blauwe gebieden tonen aan waar de nieuwe sterren geboren
worden.*

Proefschrift tot het behalen van de academische graad van
Doctor in de Wetenschappen: Sterrenkunde
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Dankwoord

Op de vooravond van mijn publieke verdediging is het ook tijd om eens te reflecteren over hoe dit werk tot stand is gekomen. Sterrenkunde heeft me van jongs af aan altijd geïnteresseerd; op de vraag wat ik later wou worden, antwoordde ik dan ook steevast “astronaut”. Deze interesse uitte zich onder andere in de regelmatige spreekbeurt over bijvoorbeeld het Zonnestelsel en een voorliefde voor science-fiction films en televisieseries. Ook had ik veel, doorgaans speculatieve, discussies over het heelal met mijn broer Simon, met wie ik een slaapkamer deelde. Dit soms wel eens ten koste van onze nachtrust. A small price to pay...

Na het middelbaar koos ik ervoor wiskunde te gaan studeren aan de Universiteit Gent. De (keuze)vakken over sterrenkunde die deze opleiding aanbood, hadden zeker ook een invloed bij deze beslissing. De vrienden uit de wiskunde- en fysica-opleiding hebben mij nog meer beïnvloed en zo kwam ik in het laatste jaar terecht bij professor De Rijcke voor mijn masterthesis. Dit onderzoek vond ik zodanig leuk dat ik het heel graag verder wou zetten tijdens een doctoraat bij dezelfde promotor, die dit ook zag zitten (en gelukkig ook geld had hiervoor). Professor De Rijcke werd gewoon Sven en ik bleef nog vier jaar langer in Gent. Niet als astronaut, maar een doctoraat over sterrenkunde is waarschijnlijk het meest realistische dat in de buurt komt.

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met name de bureaugenoten doorheen de jaren, Mina, Joeri, Annelies, Bert en Shivangee, die ervoor gezorgd hebben dat ik elke dag met plezier naar S9 kwam. Inge en Gerbrand, voor de vlotte administratieve en technische ondersteuning. Nogmaals Bert en zijn liefde voor open-source, waardoor hij het niet erg vond dat ik dezelfde L^AT_EX-template gebruikte voor deze thesis. En last but not least zou ik graag Sven enorm willen bedanken, om mij de kans te geven aan dit doctoraat te beginnen, voor de steun en nuttige discussies, voor het geloof in mij en voor de vrijheid die ik gekregen heb om mijn eigen onderzoeksinteresses te gaan ontdekken.

Robbert Verbeke
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Summary

Let us think the unthinkable, let us do the undoable, let us prepare to grapple with the ineffable itself, and see if we may not eff it after all.

Douglas Adams

LCDM is the current standard model of cosmology and has been very successful in describing the universe. For example by accurately reproducing its large-scale structure and explaining the observed primordial elemental abundances. However, Λ CDM starts encountering problems in the regime of dwarf galaxies. The shallow gravitational potentials of these low-mass systems make them more susceptible to both internal and external mechanisms acting on them. This makes dwarf galaxies ideal systems to test cosmological and galaxy evolution theories. In this work, we have used computer simulations to model dwarf galaxies of different masses to see how they fit within these cosmological models. To properly probe the regime of dwarf galaxies, it is important to include all physical processes that have an influence on their evolution. We furthermore stress the importance of analyzing these simulations in an observationally motivated way.

We have studied the effect an infalling gas cloud has on the evolution of an isolated dwarf galaxy. We were especially interested to see if this interaction could lead to a significant enhancement of the star formation rate of the galaxy. *Blue Compact Dwarfs* are such systems with high star formation rates and have properties that differ from other dwarfs. For example, they have smaller scale lengths (in other words, they are more compact) and steeply rising rotation curves. We show that the simulated galaxies that experience an increase in their star formation rate show these properties as well. This is a consequence of a deepening of the gravitational potential caused by the infalling gas cloud. Stellar feedback

following the starburst rapidly removes material from the central regions, making this effect only temporary.

As we incorporated more relevant physics in our model, such as heating by the cosmic ultra-violet background, we found that one crucial ingredient was still missing from our models to reproduce dwarf galaxies that resemble observed ones. The answer was found in the first generation of stars that formed after the big bang, whose effects had thus far not been taken into account. The star formation in the early universe produced significantly more high-mass stars, which resulted in more UV radiation and more and more powerful supernovae explosions. This intense stellar feedback has a strong effect on the present-day properties of the simulated dwarfs and was shown to be of crucial importance to produce a realistic population of dwarf galaxies. This led to the MORIA (Models of Realistic dwarfs In Action) suite of simulations.

Due to the high resolution of the MORIA simulations, they are well suited to produce realistic synthetic observations. Using commonly used astronomical software, we analyzed the kinematics of their neutral gas contents. The observationally motivated rotation curves we obtain in this way differ from the dynamical rotation curve, which is what we would expect based on how matter is distributed throughout the galaxy. For the lowest-mass MORIA dwarfs, the gas rotation velocity is consistently lower than what we would expect. This is a possible answer to the *Too Big To Fail* problem in cosmology for gas-rich dwarfs, which states that observed low-mass galaxies inhabit smaller dark matter halos than what is theoretically expected. The halo masses are inferred from the galaxies kinematics, so if the rotational velocities are lower, they will indeed seem to inhabit lower-mass halos than they actually do. This explains the apparent Too Big To Fail problem.

The common theme throughout this work is that we investigate some of the open issues Λ CDM has in the regime of small galaxies and see if they can be resolved by properly modeling and analyzing them.

Nederlandstalige samenvatting

Science progresses best when
observations force us to alter our
preconception.

Vera Rubin

ACDM is het huidige standaardmodel voor kosmologie en is heel succesvol gebleken in het beschrijven van het universum. Bijvoorbeeld door het correct reproduceren van de grootschalige structuur van het universum en het verklaren van de geobserveerde primordiale abundanties van chemische elementen. In het regime van dwergsterrenstelsels botst Λ CDM echter op wat problemen. De zwakke zwaartekrachtspotentialen van deze systemen maken hen enorm gevoelig voor zowel interne als externe mechanismen die op hen inwerken. Dit maakt dwergsterrenstelsels ideale systemen om theorieën voor kosmologie en galaxie-evolutie te testen. In deze thesis hebben we computersimulaties gebruikt om dwergsterrenstelsels met verschillende massa's te modelleren om te zien hoe deze passen binnen grotere kosmologische modellen. Om tot het regime van de dwergsterrenstelsels door te dringen, is het belangrijk om alle fysieke processen die een invloed hebben in hun evolutie in rekening te brengen. Daarbovenop benadrukken we hierbij het belang van de simulaties te analyseren op een observationeel gemotiveerde manier.

We hebben het effect dat een invallende gaswolk heeft op de evolutie van een geïsoleerd dwergsterrenstelsel bestudeerd. We waren in het bijzonder geïnteresseerd om te zien of deze interactie een significante opflakking in het stervormingstempo teweeg kon brengen. *Blue Compact Dwarfs* zijn zulke systemen met hoge stervormingstempo's en hebben eigenschappen verschillend van andere dwergsterrenstelsels. Zo hebben ze bijvoorbeeld kleinere schaallengtes (in andere woorden, ze zijn compacter) en hebben ze scherp stijgende rotatiecurves. We tonen dat de gesimuleerde sterrenstelsels die een stijging in hun stervormingstempo ondergaan deze eigenschappen ook vertonen. Dit is een gevolg van het

dieper worden van de zwaartekrachtspotentiaal, veroorzaakt door de invallende gaswolk. Supernova-explosies, die afgaan na de intense periode van stervorming, verwijderen efficiënt gas uit de centrale regio's, waardoor dit effect slecht tijdelijk is.

Naarmate we meer relevante fysica aan onze modellen toevoegen, zoals opwarming door de kosmisch ultra-violette achtergrond, bleek dat er nog steeds een cruciaal ingrediënt ontbrak in onze simulatiecode om realistische dwergsterrenstelsels te reproduceren. Het antwoord kwam in de vorm van de eerste generatie sterren die ontstonden na de Big Bang en wiens effect tot dan toe niet in rekening was gebracht. Stervorming in het vroege universum produceerde significant meer massieve sterren, met meer UV straling en meer, en bovendien krachtigere, supernova-explosies als gevolg. Deze intense stellaire feedback heeft een enorm effect op de huidige eigenschappen van de gesimuleerde dwergen en bleek van cruciaal belang om een realistische populatie van dwergsterrenstelsels te produceren. Dit leidde tot de MORIA (Models of Realistic dwarfs In Action) simulaties.

Door hun hoge resolutie zijn de MORIA simulaties goed geschikt om realistische synthetische observaties te produceren. Door gebruik te maken van observationele astronomische software kunnen we de kinematica van hun neutraal waterstof analyseren. De observationeel gemotiveerde rotatiecurve die we op deze manier bekomen verschilt echter van de dynamische rotatiecurve, die we zouden verwachten op basis van de verdeling van de materie. Voor de MORIA dwergen met de laagste massa's is de rotatiecurve van het gas consistent lager dan wat we zouden verwachten. Dit is een mogelijk antwoord op het zogenaamde *Too Big To Fail* probleem in Λ CDM voor gasrijke dwergsterrenstelsels, dat zegt dat de geobserveerde dwergen in donkere materie halo's met lagere massa's leven dan wat theoretisch voorspeld wordt. De massa's van die halo's wordt afgeleid uit de gaskinematica. Als de geobserveerde rotatiesnelheden dus te laag zijn, lijken deze sterrenstelsels in een te lage donkere materie halo te leven, wat het schijnbare *Too Big To Fail* probleem verklaart.

De rode draad doorheen dit werk is dat we enkele van de openstaande problemen die het standaard kosmologisch model heeft in het regime van kleine sterrenstelsels gaan onderzoeken en kijken of deze opgelost kunnen worden. Dit door gebruik te maken van betere modellen en realistischere analysemethoden.

Introduction

There's something that doesn't make sense. Let's go and poke it with a stick.

The Eleventh Doctor

1.1 Cosmology

1.1.1 Getting to the standard model

TRYING to understand the large scale behaviour of the universe is the area of cosmology. For the start of modern cosmology, one has to look to general relativity (Einstein, 1916), which describes gravity as a consequence of the geometry of spacetime. Assuming a homogeneous and isotropic universe, one can derive the Robertson-Walker metric

$$ds^2 = c^2 dt^2 - R(t)^2 \left(\frac{dr^2}{1 - kr^2} + r^2(d\theta^2 + \sin^2 \theta d\phi^2) \right), \quad (1.1)$$

where k denotes the curvature of the universe and $R(t)$ is its scale factor. Applying the field equations from general relativity, we get the Friedmann-Lemaître equations (Friedmann, 1922; Lemaître, 1927)

$$\frac{3}{R^2} \left(k + \frac{\dot{R}^2}{c^2} \right) = \frac{8\pi G}{c^2} \rho \quad (1.2)$$

$$\frac{3\ddot{R}}{R} + 4\pi G \left(\rho + \frac{3p}{c^2} \right) = 0. \quad (1.3)$$

These tell us how the scale factor evolves in function of the matter (or equivalently the energy) density in the universe. This raises the question what matter is present in our universe. First off, it contains radiation, which has by now been

diluted into insignificance by the expansion of the universe. Next to this, the universe also contains all the matter that we see, e.g. stars and interstellar gas and dust. This form of matter we will call baryonic matter throughout this thesis. However, the need for some unseen form of matter was soon recognized. Kapteyn (1922) inferred from the higher than expected velocities of stars that there must be some form of unseen matter. On larger scales, the need for *dunkle Materie*, or *dark matter*, was recognized from the motions of galaxies in the outskirts of the Coma cluster (Zwicky, 1933). Later, evidence for more mass than can be accounted for by the luminous matter was inferred from the rotation curves from galaxies (Babcock, 1939; Rubin & Ford, 1970; Bosma, 1978; Rubin *et al.*, 1980). All these observations show us that there is about five to six times more dark matter than baryonic matter in the universe.

Other evidence for the nature of dark matter can be derived from the Bullet cluster, a system of merging galaxy clusters. Observations have shown that the gas (heated by the collision and emitting in X-ray) and galaxies are displaced. While $\sim 90\%$ of the observed baryonic mass is in the form of the X-ray gas, weak lensing has shown that most of the matter distribution follows the galaxies (Clowe *et al.*, 2006). This can be seen in Figure 1.1. Furthermore, the displacement puts constraints on the self-interaction cross-section of dark matter. Markevitch *et al.* (2004) found that $\sigma/m < 1 \text{ cm}^2 \text{ g}^{-1}$, in agreement with collisionless (i.e. cold) dark matter.

Solving the Friedmann-Lemaître equations (1.2) and (1.3) for a universe filled with radiation and matter (both baryonic and dark), we find a scale factor that vanishes at some time in the past (called the Big Bang). From there, it can keep growing but at a diminishing rate, resulting in a universe that will exist forever. An other option is that the universe expands at first, but starts shrinking again, resulting in a so called Big Crunch. Using supernovae distance measurements, Riess *et al.* (1998) and Perlmutter *et al.* (1999) independently found that the expansion of the universe is faster now than at earlier epochs. This expansion of the universe is only possible if we have a fluid with negative pressure, acting as a form of anti-gravity. The basic assumption is that this comes from *dark energy*, which is usually associated with vacuum energy.

This model for cosmology is named Λ CDM, for its inclusion of dark energy (Λ) and cold dark matter (CDM). In this age of high precision cosmology, the cosmological parameters for Λ CDM are well determined. The Hubble constant, which tells us the current expansion rate, is $H_0 = 67.74 \pm 0.46 \text{ km s}^{-1} \text{ Mpc}^{-1}$. At the present day, the baryon density parameter is $\Omega_b = 0.0486 \pm 0.0010$, the total matter density parameter is $\Omega_m = 0.3089 \pm 0.0062$, the dark energy density parameter is $\Omega_\Lambda = 0.6911 \pm 0.0062$, and the radiation density parameter is $\Omega_\gamma = 9.12 \times 10^{-5}$.



Figure 1.1: Composite image of the Bullet Cluster. X-ray observations in red (taken from Markevitch, 2006), optical observations, and the mass distribution in blue (taken from Clowe *et al.*, 2006).

(inferred from the cosmic microwave background radiation: Planck Collaboration *et al.*, 2016a). Adding the density parameters, we find that we live in a flat universe ($\Omega_{\text{tot}} = 1$).

1.1.2 Big bang nucleosynthesis

Using the cosmological parameters obtained from the cosmic microwave background (CMB), one can predict how the formation of elements heavier than hydrogen will go shortly after the big bang. This process is called big bang nucleosynthesis. This will be affected by the material available for nucleosynthesis (i.e. Ω_b) and the temperature (or density) evolution of the universe, dependent on the different density parameters through the Friedmann-Lemaître equations (1.2) and (1.3). In the early universe, the radiation density is dominant. Predictions can be seen in Figure 1.2. The parameter η on the horizontal axis is the ratio of the number of protons in our universe and the number of photons in the

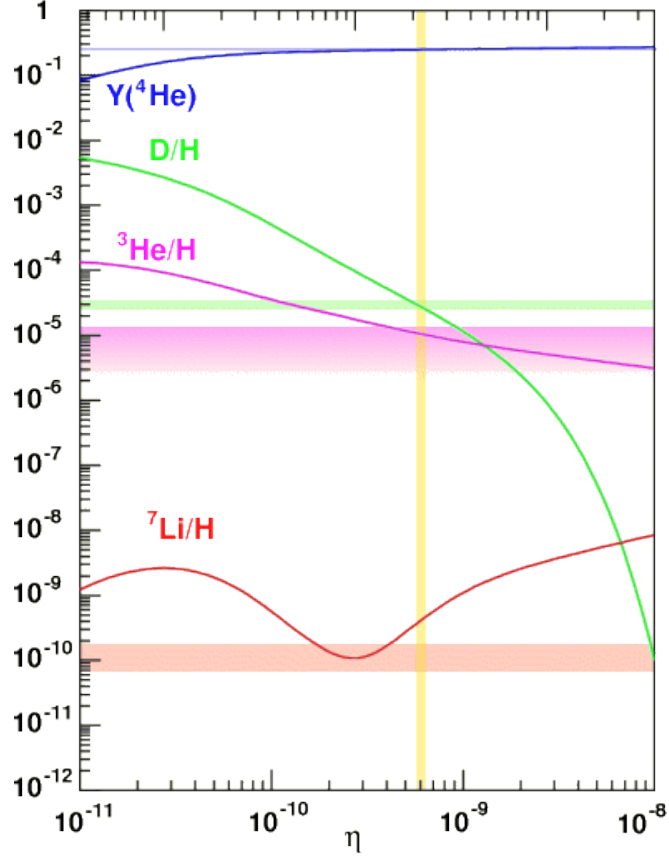


Figure 1.2: Results from big bang nucleosynthesis. See text for an in-depth explanation.

CMB:

$$\eta = \frac{n_p}{n_\gamma}. \quad (1.4)$$

The horizontal bands show the observed primordial abundance of different elements. The different curves give the primordial nucleosynthesis predictions for a given η . The yellow vertical line indicate the value of η that simultaneously reproduces all the observed abundances. The primordial ${}^7\text{Li}$ abundance is over-predicted, however this can be attributed to the uncertainty of the observed

1.1 Cosmology

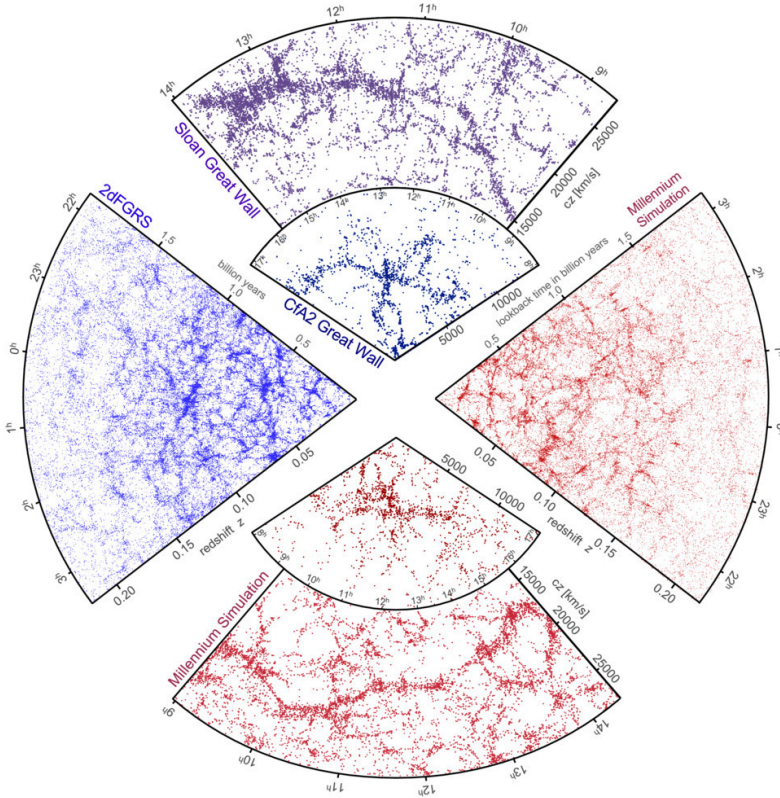


Figure 1.3: Large scale structure in the Millennium simulation (in red) compared to several observational surveys (in blue). Figure taken from Springel *et al.* (2006).

primordial abundance. The successful prediction of the elemental abundances shows that Λ CDM describes the early universe very well.

1.1.3 Cosmological simulations

The evolution of structure under gravity in an expanding universe is highly non-linear. Using approximations, one can calculate the formation redshift of structures with a certain mass (Lemaître, 1933; Tolman, 1934; Bondi, 1947). However, to properly investigate how structure formation occurs in a Λ CDM universe, we need to perform cosmological simulations. However, this can be very computa-

tionally expensive. The Millennium simulation (Springel *et al.*, 2005) used over 10^9 particles, each with a mass of $\sim 10^9 M_\odot$, to simulate a cubic region ~ 700 comoving Mpc on a side. Figure 1.3 shows the large scale structure of our universe, obtained using different surveys, in blue. Similar synthetic observations were obtained for the Millennium simulation. As can be seen, the large scale structure of the universe is well reproduced.

This and subsequent simulations only included gravity and the expansion of the universe, as described by equations (1.2) and (1.3). These simulations are then post-processed using semi-analytical methods to investigate the visible properties of galaxies. These are typically called *dark matter only (DMO)* simulations. Advances in computational resources and more efficient algorithms have led to the inclusion of hydrodynamics and astrophysical prescriptions such as star formation, stellar feedback and black hole growth in cosmological simulations. The Illustris (Vogelsberger *et al.*, 2014) and EAGLE (Evolution and Assembly of GaLaxies and their Environments: Schaye *et al.*, 2015) projects are two such *hydrodynamical simulations*. In contrast to DMO simulations, for which the physics included are well known, the physics of galaxy formation is not so well understood. Furthermore, these simulations lack the resolution to properly capture all these processes so they have to implement them using subgrid prescription (e.g. star formation is implemented as the formation of stellar particles representing a stellar population, not individual stars). Typically, the subgrid implementation is tuned so that one or more observational properties is reproduced. For the Illustris project, this is the relation between stellar mass and dark matter mass and the total amount of star formation in the universe as a function of time. For the EAGLE project on the other hand, the galaxy stellar mass was chosen. Other observational properties of galaxies can then be investigated. In this sense, hydrodynamical simulations can be used to investigate which astrophysical processes play a role in galaxy evolution and in what sense (e.g. Sijacki *et al.*, 2015; Bahé *et al.*, 2017). These hydrodynamical cosmological simulations are able to reproduce a broad range of observed galaxy properties (e.g. Sales *et al.*, 2015; Ferrero *et al.*, 2017).

1.1.4 Open issues

Given all this, Λ CDM is generally considered as the current standard model for cosmology and cosmic structure formation and is a superbly successful theory on large, super-galactic distance scales (Mamon *et al.*, 2017; Rodríguez-Puebla *et al.*, 2016; Planck Collaboration *et al.*, 2016b; Cai *et al.*, 2014; Suzuki *et al.*, 2012). However, there are some unresolved issues. Probably the biggest one is the failure so far to directly detect dark matter. Also, naive theoretical calculations of the dark energy, assuming it corresponds to vacuum energy, give a value

1.1 Cosmology

many orders of magnitude larger than observed. On top of this, when going to smaller, sub-galactic scales, and especially in the regime of dwarf galaxies, Λ CDM encounters a number of persistent problems.

Since gravity is scale free, the number of subhalos above a certain mass threshold is expected to follow a power law. This behaviour is indeed found in simulations, with hundreds of satellite galaxies around a Milky Way (MW) analogue (Klypin *et al.*, 1999; Moore *et al.*, 1999). This is an order of magnitude higher than the observed number of satellite of the MW. This is called the *Missing Satellites Problem*. It can be alleviated by assuming that not every dark matter substructure will host a galaxy. Indeed, Sawala *et al.* (2015) showed that reionization prevents galaxies to form when $M_{\text{halo}} \lesssim 3 \times 10^9 M_{\odot}$. The accretion histories of the halos and environmental effects also play an important role in populating halos with galaxies (Sawala *et al.*, 2016b).

There is disagreement on the inner density profiles of dark matter halos look like. DMO simulations predict a *cusp*, i.e. the density diverges towards the center (Dubinski & Carlberg, 1991). More specifically, the density is given by the Navarro-Frenk-White profile (NFW; Navarro *et al.*, 1996b)

$$\rho_{\text{DM}}(r) = \frac{\rho_s}{\left(1 + \frac{r}{r_s}\right)^2}, \quad (1.5)$$

with r_s and ρ_s the characteristic radius and density, respectively. This profile holds for all halo masses. While for massive galaxies dark matter only dominates in the outskirts, dwarf galaxies are dark matter dominated everywhere. This makes them ideal candidates to test whether their dark matter halos indeed have a cusp. Observations of dwarf galaxy rotation curves have shown that they in fact have a core, i.e. a flat density profile (de Blok *et al.*, 2008; Salucci *et al.*, 2012). However, including baryonic effects in the simulations might change the dark matter density in the inner region. Indeed, for example, gas blown out because of supernova feedback might dynamically heat the dark matter halo, lowering the inner densities (e.g. Pontzen & Governato, 2012; Cloet-Osselaer *et al.*, 2012; Read *et al.*, 2016). Di Cintio *et al.* (2014b) found that the efficiency of core formation is dependent on the star formation efficiency $M_{\star}/M_{\text{halo}}$: it is most efficient for $M_{\star}/M_{\text{halo}} \approx 0.005$, or using abundance matching (Moster *et al.*, 2013), $M_{\text{halo}} \approx 6 \times 10^{10} M_{\odot}$ and $M_{\star} \approx 3 \times 10^8 M_{\odot}$. For lower values, the total energy from supernovae results in less efficient core formation. For higher values, the galaxies become too centrally concentrated and less efficient in driving the large outflows necessary for core formation. Chan *et al.* (2015) find similar results, but with a slightly different dependence of core formation on $M_{\star}/M_{\text{halo}}$. Secondary effects, such as the star formation history of a galaxy (Oñorbe *et al.*, 2015), can

influence the efficiency of core formation.

Another problem is called *Too Big Too Fail*, or TBTF, first formulated in the context of the Local Group and in a sense an extension of the missing satellites problem. Given the many factors that suppress star formation in dwarf galaxies, such as supernova feedback and the cosmic UV background, visible dwarf galaxies are expected to reside in relatively scarce high- v_{circ} dark-matter halos. This would also agree with their small observed number density. However, most observed MW satellites have circular velocities $v_{\text{circ}} < 30 \text{ km s}^{-1}$, estimated from their stellar kinematics, indicating that these satellites seem to live in low- v_{circ} subhalos, which are too abundant in comparison with the observed number of MW satellites (Boylan-Kolchin *et al.*, 2011, 2012). The TBTF problem is also present for the satellite system of Andromeda (Tollerud *et al.*, 2014) and for field dwarfs in the Local Group and Local Volume (e.g. Ferrero *et al.*, 2012; Garrison-Kimmel *et al.*, 2014; Papastergis *et al.*, 2015).

Several possible solutions to this problem have been suggested. For example, if the MW were to have a smaller virial mass, then it would also host a smaller number of massive subhalos (Wang *et al.*, 2012). Another way out is to take into account the fact that baryonic processes, such as supernova feedback, can flatten the inner dark-matter density distribution, converting a high- v_{circ} cuspy density profile into a low- v_{circ} cored one at constant halo mass. By fitting the mass-dependent DC14 profile (Di Cintio *et al.*, 2014a) to the kinematical data of the Local Group dwarf galaxies, Brook & Di Cintio (2015a) found that dwarf galaxies inhabit more massive halos than previously thought, thus alleviating the TBTF problem. Other effects that help reduce dwarf galaxy circular velocities in the context of the Local Group include tidal stripping (Sawala *et al.*, 2016b).

In their analysis of the TBTF problem in field dwarfs, where only internal baryonic effects can be invoked to reduce halo circular velocities, (Papastergis & Shankar, 2016, henceforth referred to as P16) use abundance matching to derive the relation between the observed H I rotation velocity inferred from the galaxy 21cm emission line profile, W_{50} , and the maximum halo circular velocity $v_{h,\text{max}}$ such that the halo velocity function (VF) found in simulations (Sawala *et al.*, 2015) corresponds to the observed field galaxy VF (Haynes *et al.*, 2011; Klypin *et al.*, 2015). Hereafter, we will refer to this relation between W_{50} and $v_{h,\text{max}}$ as the P16-relation. Then, these authors fit NFW (Navarro *et al.*, 1996b) and DC14 profiles to the outer-most datapoint of the rotation curves of a set of field dwarf galaxies to infer their $v_{h,\text{max}}$. This allows them to put individual $v_{\text{rot,HI}} - v_{h,\text{max}}$ datapoints on the inferred statistical relation. As these authors note: “ Λ CDM can be considered successful only if the position of individual galaxies on the $W_{50} - v_{h,\text{max}}$ plane is consistent with the relation needed to reproduce the measured VF of galaxies.” As it turns out, the individual galaxies are not consistent

1.1 Cosmology

with the expected P16-relation.

The discrepancy between these results and those from Brook & Di Cintio (2015a) results from the radius at which the circular velocity is measured: for measurements beyond the core radius ($\gtrsim 2\text{ kpc}$), fitting a DC14 profile gives similar results as using a cusped NFW profile. The TBTF problem can therefore not be (fully) explained by core creation alone (see also Papastergis & Ponomareva, 2017). The TBTF problem will be the focus of Chapter 5.

The 11 classical MW satellites were found to all lie in a highly flattened planar structure (Lynden-Bell, 1976; Kroupa *et al.*, 2005). Nine of them have their angular momentum vector perpendicular to this plane of satellites (in other words, they are rotating within the plane), eight of which are co-rotating (Metz *et al.*, 2009b; Pawlowski & Kroupa, 2013), indicating that this *Vast Polar Structure* (VPOS) is not transient. The VPOS also consists of fainter satellites (Metz *et al.*, 2009a), as well as globular clusters and stellar and gaseous streams (Pawlowski *et al.*, 2012). Similar planes have been found around e.g. M31 (Great Plane of Andromeda, GPoA; Ibata *et al.*, 2013), NGC 3109 (Bellazzini *et al.*, 2013), and Centaurus A (Tully *et al.*, 2015). On the other hand, Λ CDM predicts an isotropic distribution for satellite systems, making identifying a structure like the VPOS or GPoA very unlikely in cosmological simulations (e.g. Pawlowski *et al.*, 2014; Ibata *et al.*, 2014). However, Sawala *et al.* (2016a) find satellite planes around the centrals in their 12 Local Group simulations that are as flattened the VPOS and GPoA. They attribute this to the fact that satellites fall in to a group along filaments in the cosmic web (Libeskind *et al.*, 2005). This claim is contested by Pawlowski *et al.* (2015), who point out that their way of measuring the flattening of the plane might be flawed. Clearly, this issue is far from settled.

Dwarf galaxies can form in the tidal arms of interacting spiral galaxies (Mirabel *et al.*, 1992). Since such tidal dwarf galaxies (TDGs) form out of dynamically cold gas, they should not be able to capture any of the dynamically warm dark matter from the halo and as such, should have mass to light ratios close to unity (Barnes & Hernquist, 1992). Such seemingly DM-free objects have been observed (Lelli *et al.*, 2015; Sweet *et al.*, 2016), although it is not clear whether these systems are in dynamical equilibrium and consequently, if their dynamically inferred masses can be trusted. On the other hand, the TDGs around NGC 5291 were found to have a significant dark component (Bournaud *et al.*, 2007). Bournaud & Duc (2006) performed self-consistent simulations of TDGs and found $\sim 25\%$ of them to be long-lived. Furthermore, they found that TDGs should constitute a small but significant fraction of the overall dwarf galaxy population. This constitutes a problem for Λ CDM, as one would expect to observe both DM-free and DM-dominated dwarf galaxies.

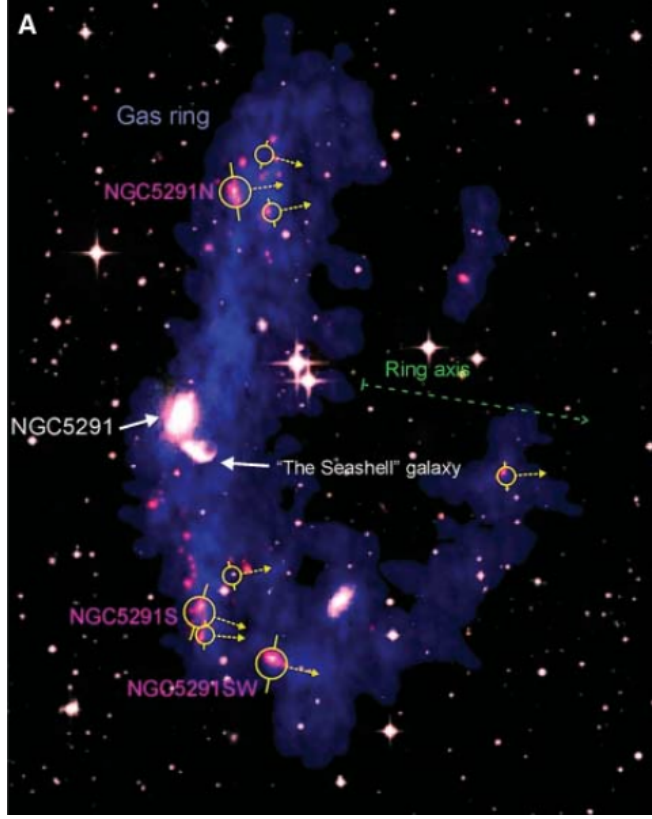


Figure 1.4: Tidal dwarf galaxies forming in the tidal tails around the interacting galaxies NGC 5291 and the Seashell galaxy. Figure taken from Bournaud *et al.* (2007).

However, we argue that including the effects of the cosmic UV background (which the simulations of Bournaud & Duc (2006) do not take into account), might have a serious effect on the number of formed TDG and on their survivability. To the best of our knowledge, full cosmological simulations (e.g. Schaye *et al.*, 2015; Vogelsberger *et al.*, 2014) do not predict a significant fraction of old TDGs at $z = 0$. But as far as we know, there has been no study undertaken into the formation and evolution of TDGs within large cosmological simulations.

Λ CDM currently has some other open issues as well, such as a discrepancy in

1.1 Cosmology

the exact value of the Hubble constant (Suyu *et al.*, 2014), the missing baryon problem (Persic & Salucci, 1992; Bell *et al.*, 2003), or predictions for the primordial Lithium abundance (Korn *et al.*, 2006; Piau *et al.*, 2006). We do not go into more detail on this.

1.1.5 Alternatives

While a lot of researchers see these issues as evidence that the models for e.g. galaxy formation are missing some crucial ingredient(s), others see it as fundamental for Λ CDM.

As a first alternative for Λ CDM, one can step away from the assumption that dark matter has to be cold and collisionless. Warm dark matter (WDM), which erases the smallest density fluctuations due to free streaming, and self-interacting dark matter (SIDM), which has a non-zero cross-section, are possible alternatives. e.g. Zavala *et al.* (2009) and Vogelsberger *et al.* (2016) for simulations in such Λ WMD and Λ SIDM cosmologies.

More drastically, an alternative to the dark matter hypothesis as a solution for the “missing mass problem” was proposed by Milgrom (1983), stating that in the limit of low acceleration (e.g. the outskirts of spiral galaxies), Newtonian dynamics (and by extension General Relativity) is no longer valid. Modified Newtonian Dynamics (MoND) states that there exists a certain acceleration scale a_0 , below which the dynamics increasingly differ than those predicted by Newton. In other words, the acceleration a an object with mass m experience due to the gravitational force F , is now given by

$$F = m\mu\left(\frac{a}{a_0}\right)a, \quad (1.6)$$

where F is the same (Newtonian) gravitational force and $\mu(x)$ is an “interpolating function”. To recover classical Newtonian dynamics for $a \gg a_0$, we need $\mu(x) \rightarrow 1$ for $x \gg 1$. If $\mu(x) \rightarrow x$ for $x \ll 1$, then we also recover flat rotation curves.

MoND is able to explain the kinematics of galaxies very well. By construction it can reproduce the observed flat rotation curves of galaxies. On top of this, the baryonic Tully-Fisher relation (BTFR), which relates the total baryonic mass of a galaxy to the value of the flat part of the rotation curve (v_f), follows naturally from MoND. As an extension to this, McGaugh *et al.* (2016) found that the circular velocity at any radius can be derived solely from the distribution of the baryonic matter, by

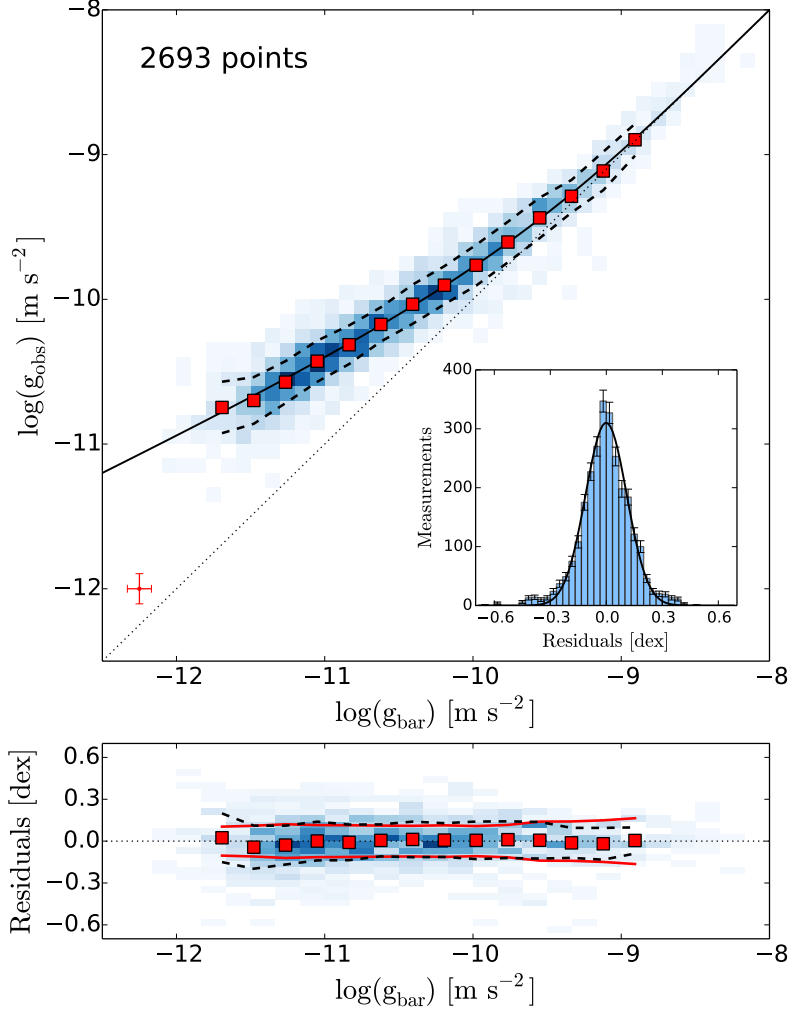


Figure 1.5: Radial acceleration relation for late-type galaxies. Figure taken from McGaugh *et al.* (2016).

$$g_{\text{obs}} = \frac{g_{\text{bar}}}{1 - e^{-\sqrt{g_{\text{bar}}/g_0}}}, \quad (1.7)$$

where g_{bar} is the expected Newtonian acceleration due to the baryons. By first

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(numerically) solving the Poisson equation

$$\nabla^2 \Phi_{\text{bar}} = 4\pi G \rho_{\text{bar}}, \quad (1.8)$$

one can obtain the Newtonian baryonic acceleration:

$$g_{\text{bar}} = \left| \frac{\partial \Phi_{\text{bar}}}{\partial R} \right|. \quad (1.9)$$

Their results are shown in Figure 1.5. This *Radial Acceleration Relation* (RAR) holds for all galaxies, irrespective of their mass, size or type (Lelli *et al.*, 2017b). It has exactly the behaviour necessary for the interpolation function of MoND. They found $g_0 = 1.20 \times 10^{-10} \text{ m s}^{-2}$.

The BTFR and RAR are easily explained in MoND. In any dark matter theory on the other hand, this would require a very tight but non-linear relation between the baryons and dark matter, which is not straightforward. However, galaxy simulations within Λ CDM have been able to reproduce the BTFR (e.g. Verbeke *et al.*, 2015) and even the RAR (Keller & Wadsley, 2017), although not all simulations arrive at the observed g_0 (Ludlow *et al.*, 2017). On top of this, within MoND, the planes of satellites are easily explained as a result of these satellites being formed during an interaction between the two massive galaxies. These TDGs would form in a coherent structure and, because of MoND, have mass-to-light ratios higher than unity.

This does not mean MoND can explain everything perfectly. While it works very well for galaxies, it encounters some problems on larger scales. For example, it suffers from a missing mass problem for galaxy clusters (The & White, 1988) and the Bullet Cluster is not easily explained in MoND. See Dodelson (2011), and references therein, for an overview of the issues MoND is facing. As discussed above, Λ CDM was designed to describe the large scales of the universe, but encounters problems for the lowest-mass galaxies. In the same sense, it is not surprising that MoND, which was designed to work well on galaxy-scales, has trouble with clusters and even larger scales.

An other way of looking at the fundamental differences between MoND and Λ CDM, is that the former looks for an answer to the missing mass problem in an adjustment to General Relativity as a description for gravity, while the latter looks for physics beyond the Standard Model (of particle physics), since no known particle is a suitable candidate for dark matter. In this sense, the solution might lie within a theory for quantum gravity. For example, *emergent gravity* (Verlinde, 2016) predicts a “dark gravity force”, reminiscent of MoND for galaxies. On top of this, it has no missing mass problem for clusters, in contrast with MoND.

However, Lelli *et al.* (2017a) found emergent gravity to be inconsistent with the RAR. But the idea that the missing mass problem can be solved within the framework of quantum gravity theories is definitely promising.

1.2 Dwarf galaxies

As shown earlier, Λ CDM suffers from some problems at the dwarf galaxy scale. Furthermore, these low-mass systems are the most numerous types of galaxies in the universe and formed the earliest. Their shallow gravitational potentials make them ideal testbeds for testing galaxy evolution models. For this reason, the work done for this PhD had focused strongly on dwarf galaxies and, as such, we now give a brief introduction on these types of objects.

Dwarf galaxies are small compared to elliptical or spiral galaxies. The most massive ones have stellar masses of $\sim 10^9 M_\odot$ and a V-band magnitude of $M_V \sim -17$ mag. They typically have larger sizes and lower surface brightnesses than globular clusters (GCs) at a given luminosity. They can be divided in several classes (Mateo, 1998; Tolstoy *et al.*, 2009).

- **Early-types.** Dwarf ellipticals (dEs) and dwarf spheroidals (dSphs) are gas-poor and thus have very little active star formation (SF). The difference between dEs and dSphs is mostly based on their magnitudes (and thus their mass). dSphs have $M_B \gtrsim -14$ mag, while dEs have $-14 \text{ mag} \gtrsim M_B \gtrsim -18$ mag, so one might argue that dSphs are fainter versions of dEs. Their name is derived from the fact that they have elliptical or spheroidal isophotes. dEs and dSphs typically have little to no on-going star formation and are mostly not or slowly rotating. However, see e.g. Toloba *et al.* (2015) for cases of dEs with significant rotation. These type of dwarfs are mostly found in dense environments like galaxy clusters. Ultra-faint dwarfs (UFDs) can be seen as another extension of these types of objects to even lower luminosities.
- **Late-types.** Dwarf irregulars (dIrrs) on the other hand, are irregular, in the sense that their isophotes deviate strongly from ellipses. They have significant gas fractions and active star formation. They generally have coherent rotation in their gas component. This is however not necessarily true for their stellar component, which seem to be mostly pressure supported, just like dEs/dSphs (Kirby *et al.*, 2014; Koleva *et al.*, 2014; Wheeler *et al.*, 2017) dIrrs are mostly found in less dense environments. This off-set between the location where dEs and dIrrs are found, is called the morphology-density relationship. Blue compact dwarfs (BCDs) are dIrrs that are currently going through an intense period of star formation, known as a starburst. BCDs

1.2 Dwarf galaxies

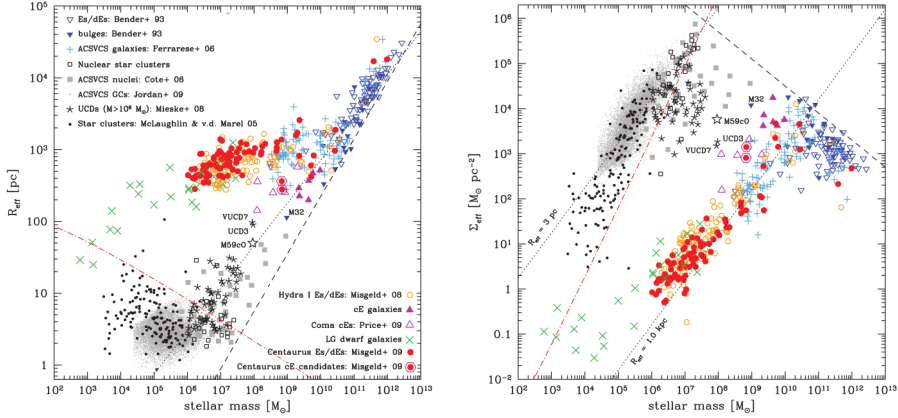


Figure 1.6: Half-light radius (left) and central surface brightness (right) of different types of stellar systems. Figure taken from Misgeld & Hilker (2011).

and the trigger mechanism of their starburst is the subject of Chapter 3, in which we give a more extensive overview of their properties.

- **Transition types.** Some dwarf galaxies are classified in an intermediary class, the transition type dwarfs (dT). They have the same characteristics as dSphs, but also have gas present.
- **Ultra compact dwarfs.** UCDs are much more compact than dSphs of similar luminosity and are mostly found in galaxy clusters. They have spectra similar to other types of dwarfs and are more luminous than GCs. However, based on their structural parameters, they look more like GCs and galaxy nuclei. Based on this, it has been suggested that they are bright GCs or the remains of stripped early-type dwarf galaxies (see e.g. Evstigneeva *et al.*, 2008).

What the link is between the different types is not exactly clear. The existence of the dT class and the fact that they both have similar properties (e.g. exponential surface brightness profiles) has been seen as evidence for an evolutionary path between early-types and late-types (Grebel *et al.*, 2003). Since early-types clearly have formed stars at some point, they must have been gas-rich before and somehow lost their gas. Both internal and external mechanisms have been proposed to explain this gas-loss.

A first internal mechanism in which a galaxy might lose its gas is by converting all of its gas into stars. However, supernovae explosions following star formation will heat and disperse its surrounding gas, locally and temporarily shutting down star formation. Given the low gravitational wells of dwarf galaxies, this temporary quenching by supernovae is very efficient, explaining the over-all lower star formation efficiency in lower-mass galaxies. This naturally brings us to a second internal gas-removal mechanism: stellar feedback can blow the gas out of the galaxy, making it unbound and unavailable for future star formation.

A gas-rich, rotationally supported dwarf galaxy orbiting around a more massive companion can be transformed in a gas-poor, pressure supported dwarf in a process called *tidal stirring* (Mayer *et al.*, 2001). Furthermore, the low-density intergalactic medium (IGM) can have a significant influence on a galaxy and possibly completely remove its gas content through *ram-pressure stripping* (Gunn & Gott, 1972; Mayer *et al.*, 2006). The cosmic UV background will also have a strong influence on how much gas a galaxy can retain (Efstathiou, 1992).

Given the morphology density relation, which states that early-type dwarfs are more likely found in dense environments like groups or clusters and late-type dwarfs more in isolation, the external effects seem to be the dominant mechanisms in gas removal and thus a transition from dIrrs to dEs/dSphs, possibly aided by internal mechanisms.

1.3 This work

For this PhD, we try to answer some of the open questions regarding low-mass galaxies within the Λ CDM paradigm. Work done in the Dwarf Galaxy research group at Ghent University includes observational studies of dwarf galaxies (e.g. De Rijcke *et al.*, 2007; de Rijcke *et al.*, 2009; Koleva *et al.*, 2009a,b), as well as efforts to model them in computer simulations (e.g. Valcke *et al.*, 2008; Schroyen *et al.*, 2011; Cloet-Osselaer *et al.*, 2014; Vandenbroucke *et al.*, 2016). The research presented in this thesis falls within the latter category, but with a strong emphasis on comparing the simulated galaxies with nature in an observationally motivated way. While I was able to build on the efforts of my predecessors, I have run and analyzed the simulations myself. In Chapter 2, we discuss the techniques used for the galaxy simulations. Chapter 3 is based on Verbeke *et al.* (2014) and discusses the formation of BCDs. We present a possible trigger mechanism for the starburst, an explanation for the specific properties of observed BCDs and for the different subclasses. In Chapter 4 (Verbeke *et al.*, 2015), we argue for the need of strong feedback from the first generation of stars to simulate realistic low-mass galaxies. This is done by assuming a top-heavy initial mass function

1.3 This work

for the first stars. In Chapter 5 (Verbeke *et al.*, 2017), we present synthetic H I observations of the simulated dwarf galaxies. We derive rotation curves from the H I kinematics and discuss these in the context of the Too-Big-To-Fail problem for isolated late-type dwarfs. Chapter 6 gives an overview of some of our on-going research.

Simulations

R2-D2, you know better than to
trust a strange computer!

C-3PO

TO investigate some of the open questions regarding cosmology and galaxy evolution, we make use of computer simulations. For this, we need to make a trade-off between making the models as realistic as possible and making them computationally efficient. As such, resolution limits are put in place. Given that the physics governing the evolution of galaxies span many different size scales, we are forced to implement some important processes with physical scales smaller than our resolution in a *subgrid* fashion.

Assuming a Λ CDM-cosmology, galaxies consist of dark matter, gas (mostly hydrogen) and stars. Each fluid is modelled in a different way. Dark matter is assumed to be a collisionless fluid and is thus only affected by gravity. Gas is a collisional fluid, so it is affected by the surrounding gas following the hydrodynamical equations. Furthermore, gas can cool and be heated and can form stars. Stars are collisionless as well, but they affect the gas by pumping energy and metals into their surroundings. This in turn has an affect on the way gas cools and is heated and thus how stars will form later. Modelling all these processes in an effecient way is not a trivial task.

We give an overview of the techniques we to models all these astrophysical processes. Unless stated otherwise, this work was done by members of the group before me (Valcke *et al.*, 2008; Schroyen *et al.*, 2011; Cloet-Osselaer *et al.*, 2012; Schroyen *et al.*, 2013; Vandenbroucke *et al.*, 2013; De Rijcke *et al.*, 2013; Cloet-Osselaer *et al.*, 2014).

2.1 Gravity

The acceleration by a particle i due to gravity is given by

$$\ddot{\mathbf{x}}_i = G \sum_{j \neq i}^N \frac{m_j (\mathbf{x}_i - \mathbf{x}_j)}{r_{ij}^3}, \quad (2.1)$$

where G is the gravitational constant, m_j the mass of particle j , N the number of particles and r_{ij} the distance between particle i and j . Computationally, it can become a problem when two particles get too close together as the denominator in (2.1) will become too small and thus $\ddot{\mathbf{x}}_i$ will blow up. This can be countered by adding a softening term ϵ to the equations:

$$\ddot{\mathbf{x}}_i = G \sum_{j \neq i}^N \frac{m_j (\mathbf{x}_i - \mathbf{x}_j)}{(\epsilon^2 + r_{ij}^2)^{3/2}}. \quad (2.2)$$

Introducing this softening term has negligible effect if $r_{ij} \gg \epsilon$, in which case (2.2) reduces to (2.1). However, when $r_{ij} \lesssim \epsilon$, the gravitational force is significantly reduced and vanishes for $r_{ij} = 0$. We have thus introduced a gravitational force scale in our models. In Chapter 3, we use $\epsilon = 30$ pc, while for later chapters, the softening is determined for each simulation differently, depending on its resolution.

Looking at (2.1), it is clear that if we would directly compute the gravitational force on each particle, we would need of the order $\mathcal{O}(N^2)$ operations. This is not feasible for large particle numbers (for us, typically $N \sim 10^6$), so we need a more efficient algorithm. For this, we construct a hierarchical tree (Barnes & Hut, 1986). The idea behind this tree method is that the force on a particle will be mostly contributed by the particles closest to it. First of all, a cube that contains all particles is constructed. This is the first *node*. This node is then subdivided in eight equal-sized smaller nodes. Those containing no particles get discarded and for those with exactly one particle, the construction process is finished. These nodes are then called *leaves*. Those that do contain multiple particles, get subdivided again. This is done until all particles are in a leaf. During the construction of the tree, the total mass and the center of mass location of each node is calculated and stored. After the tree has been constructed, we can calculate the gravitational force on each particle by working top-down through the tree. At each level, we compare the width of a node w to the distance r between the particle and the node's center of mass. If $w/r < \delta$, with δ a certain specified tolerance, the force between the particle and all the particles in the node is calculated by assuming that all the mass of those particles is concentrated in the center of mass of the node and we do not need to go deeper down the tree.

2.2 Hydrodynamics

If however $w/r > \delta$, we go to the next level and repeat the above process. We note that this tree has to be reconstructed for every time step. Using this tree method, the number of operations needed to compute the gravitational force has become $\mathcal{O}(N \log(N))$ rather than $\mathcal{O}(N^2)$.

Now that we have computed the forces acting on each particle, we can do the actual integration. This is done using the leapfrog method, which places the positions x and velocities v on a time staggered grid. In its most basic form, it is given by

$$\begin{aligned}\mathbf{x}_{n+1/2} &= \mathbf{x}_{n-1/2} + h\mathbf{v}_n \\ \mathbf{v}_{n+1} &= \mathbf{v}_n + h\ddot{\mathbf{x}}_{n+1/2}\end{aligned}\tag{2.3}$$

with h the step size. This scheme has several advantages. It is rather simple in implementation while still being second-order accurate. Due to the underlying structure of the leapfrog method, it conserves global properties of Hamiltonian systems like the N -body problem. For example, it maintains the time reversibility of the system. Because of this, the method implicitly also handles conservation of energy very well.

2.2 Hydrodynamics

For the gas in our simulations, which is a collisional fluid, the expression for the acceleration becomes

$$\ddot{\mathbf{x}}_i = -\frac{1}{\rho_i}\nabla p_i - G \sum_{j \neq i}^N \frac{m_j(\mathbf{x}_i - \mathbf{x}_j)}{(\epsilon^2 + r_{ij}^2)^{3/2}},\tag{2.4}$$

already adding the softening term. For the second term in (2.4), we can use the tree method described above. For the pressure term, several techniques are available. One popular method called Adaptive Mesh Refinement (AMR; e.g. Teyssier, 2002) uses a Eulerian grid that can be adaptively refined in regions where higher resolutions is required. The downside of methods with a static grid is that it is not Galilean invariant, i.e. the bulk velocity of the fluid with respect to the grid may change the solution. Another method is called Smoothed Particle Hydrodynamics (SPH; Lucy, 1977; Gingold & Monaghan, 1977; Price & Monaghan, 2004), which follows the fluid using tracer particles. By construction, SPH is Galilean invariant, but it can not handle contact discontinuities as well as AMR (Agertz *et al.*, 2007). Other methods exist that try to capture the advantage of grid-based methods while still being Galilean invariant, such as co-moving

mesh methods (Springel, 2010; Vandenbroucke & De Rijcke, 2016) or mesh-free methods (Hopkins, 2015). Throughout this thesis, we will use the SPH formalism.

The basic idea behind SPH is to treat the gas particles as smeared-out density distributions, rather than point masses. We define the contribution to the density at a point $\mathbf{x}$ arising from a particle j at position $\mathbf{x}_j$ as

$$\rho^{(j)}(\mathbf{x}) = m_j W(|\mathbf{x} - \mathbf{x}_j|, h), \quad (2.5)$$

with h now the smoothing length and $W(r, h)$ the smoothing kernel that describes the mass distribution around an SPH particle. Several choices for this kernel can be used, for example, the *cubic spline* given by

$$W(r, h) = \frac{8}{\pi h^3} \begin{cases} 1 - 6(\frac{r}{h})^2 + 6(\frac{r}{h})^3 & \text{if } 0 \leq \frac{r}{h} \leq \frac{1}{2}, \\ 2(1 - \frac{r}{h})^3 & \text{if } \frac{1}{2} \leq \frac{r}{h} \leq 1, \\ 0 & \text{if } \frac{r}{h} > 1. \end{cases} \quad (2.6)$$

The physical density at a given location $\mathbf{x}$ can then be obtained by summing the overlapping contributions from all particles,

$$\rho(\mathbf{x}) = \sum_{j=1}^N \rho^{(j)}(\mathbf{x}) = \sum_{j=1}^N m_j W(|\mathbf{x} - \mathbf{x}_j|, h). \quad (2.7)$$

It is not necessary to take a fixed smoothing length h for each particle. For each particle j , we chose a smoothing length h_j so that its kernel volume contains a constant mass for the estimated density. In other words

$$\frac{4\pi}{3} h_j^3 \rho_j = N_{\text{SPH}} \bar{m} \quad (2.8)$$

with N_{SPH} the typical number of smoothing neighbours and $\bar{m}$ an average particle mass.

All physical quantities of the fluid can now be approximated by interpolation between the known particle properties. Taking the derivative of such a quantity then simplifies to taking the derivative of the kernel. For example, the derivative of the pressure, which we need for (2.4)

$$\nabla p_i = \sum_{j=1}^N m_j \frac{p_j}{\rho_j} \nabla W(|\mathbf{x}_i - \mathbf{x}_j|, h). \quad (2.9)$$

When using variable smoothing lengths, it is recommended to symmetrize the above equation:

2.3 Subgrid physics

$$\nabla p_i = \sum_{j=1}^N m_j \left[f_i \frac{p_i}{\rho_i^2} \nabla W(|\mathbf{x}_i - \mathbf{x}_j|, h_i) + f_j \frac{p_j}{\rho_j^2} \nabla W(|\mathbf{x}_i - \mathbf{x}_j|, h_j) \right], \quad (2.10)$$

with

$$f_i = \left(1 + \frac{h_i}{3\rho_i} \frac{\partial \rho_i}{\partial h_i} \right)^{-1}. \quad (2.11)$$

The particle pressures are given by $p_i = A_i \rho_i^\gamma$, with A_i the entropy of the particle and γ the adiabatic index, which for an ideal gas is equal to 5/3. Once we have an approximation of the force acting on each SPH particle, the leapfrog method can be used to evolve the system through time.

The GADGET-2 code is publicly available and implemented the methods for gravity and hydrodynamics briefly discussed here. We refer to Springel (2005) for a more detailed description of the methods and implementation. It is known that the standard SPH prescription suffers from several numerical issues (see e.g. Springel, 2010; Hopkins, 2015, and references therein). These are most notable in processes such as ram-pressure stripping. For our simulated galaxies, which do not experience processes where the standard SPH prescription gives significantly erroneous results, corrections to the hydrodynamical scheme will be unlikely to have a significant influence on the results.

2.3 Subgrid physics

GADGET-2 was extended to include several processes important in galaxy evolution¹ (Valcke *et al.*, 2008; De Rijcke *et al.*, 2013; Vandenbroucke *et al.*, 2013).

2.3.1 Radiative cooling

The interstellar medium (ISM) predominantly constitutes of Hydrogen and Helium, but also contains heavier elements and molecules. The individual atomic elements can be neutral or (partly or fully) ionized. When an ionized element recombines with an electron, it will send out a photon corresponding to the recombination energy. This photon can escape the galaxy, resulting in an energy loss in the system. For wavelengths for which the ISM is optically thin, this escape is straight-forward. However, the ISM is optically thick for e.g. Ly α and H α . In this case, cooling can be achieved through diffusion and eventual escape

¹<https://github.ugent.be/dwarf-galaxy-group/Gadget2>

from the edge of the galaxy. In a similar way, an element transitioning from an excited state to a lower-energy level will also provide a mechanism for the ISM to cool.

In the case of high temperatures, the ISM can (locally) be fully ionized and cooling due to recombination and deexcitation will become negligible. Here, charged particles being accelerated by ions will emit thermal bremsstrahlung, allowing for the ionized ISM to cool efficiently as well.

The cooling rate of a local gas parcel will thus depend on its ionization state and chemical constitution. We include two sources of metal-enrichment in the simulations (supernovae types Ia and II), so following only two elements (Fe and Mg) allows us to infer the full chemical composition. In the absence of external heating sources, we simply have to add temperature as a third parameter to calculate the cooling rate. This is done using the CHIANTI spectral library (Dere *et al.*, 2009). To avoid having to calculate this for every gas particle on each time step, these cooling rates were tabulated (down to 10 K) and the approximate cooling rate can be obtained by a simple interpolation (De Rijcke *et al.*, 2013).

2.3.2 Advanced gas physics

For a monatomic ideal gas, the thermal energy of a gas can be written as

$$u = \frac{3}{2} \frac{nkT}{\rho}, \quad (2.12)$$

with ρ its density, n its number density, T its temperature and k the Boltzmann constant. However, in the case of a multicomponent, multiphase gas, this becomes

$$u = \frac{3}{2\rho} \left(\sum_X n_X + n_e(T) \right) kT + \sum_X \left[\sum_i n_{X,i} \left(\sum_{j=i}^i \varepsilon_{X,j} \right) \right], \quad (2.13)$$

where n_X is the number density of element X and n_e is the number density of electrons. The latter depends on the ionization state of the different elements. $n_{X,i}$ is the number density of the element X which is i times ionized and $\varepsilon_{X,j}$ is the difference in energy between the $(j-1)$ th and j th ionization state of element X .

The ideal gas law for monatomic gas relates the pressure to its number density and temperature via

$$P = nkT. \quad (2.14)$$

When the gas consists of different elements of different ionization states, this becomes

2.3 Subgrid physics

$$P = \left(\sum_X n_X + n_e(T) \right) kT := \frac{\rho kT}{\mu(T)}, \quad (2.15)$$

where the $\mu(T)$, the mean constituent mass was introduced. In the hydrodynamical scheme, we thus need this value for each gas particle to compute its pressure from its density and temperature. $\mu(T)$ depends on the chemical composition of the gas parcel and the ionization state and as such, can be calculated from its temperature and two metallicity tracers. Similar to the cooling tables, three-dimensional tables with the mean constituent mass were constructed on which can be interpolated to get an approximate value for each gas particle at each timestep.

To implement the ionization potential of the gas, the energy is split up in a kinetic part and a potential energy part

$$u = u_{\text{kin}} - \frac{\chi_{\text{H}}}{m_{\text{H}}} x, \quad (2.16)$$

where χ_{H} is the energy needed to ionize H I, m_{H} the mass of a hydrogen atom and x the ionization fraction of the gas ($x = 0$ indicating completely neutral). The evolution of the kinetic energy can then be described by

$$\frac{du_{\text{kin}}}{dt} = \frac{1}{1 + X_{\text{ukin}}} \frac{du}{dt}, \quad (2.17)$$

where we introduced the dimensionless quantity $X_{\text{ukin}} := \frac{\chi_{\text{H}}}{m_{\text{H}}} \frac{dx}{du_{\text{kin}}}$, which describes the absorption of thermal energy by the ionization potential energy reservoir. X_{ukin} is tabulated as well to increase the efficiency of the simulation code.

The effects including these advanced gas dynamics is shown in Vandenbroucke *et al.* (2013) for Sedov-Taylor blast wave. They show that a significant fraction of the energy of the blast wave is absorbed by the ionization potential. This process can not be neglected in simulations of galaxies, where supernovae explosions play an important role.

2.3.3 Ultraviolet background

The cosmic ultraviolet background (UVB; Efstathiou, 1992, and references therein) is responsible for the reionization of the universe. Next to keeping the intergalactic medium fully ionized, it also has an effect on galaxies. Indeed, galaxy formation becomes increasingly more inefficient in lower-mass halos. Reionization is thought to start at $z \sim 10.5$, with the strength of the UVB increasing until $z \sim 2$ (Faucher-Giguère *et al.*, 2009), coinciding with the peak of cosmic star formation density.

For dense neutral media, all ionizing photons will be absorbed by the H I, making the UVB unable to penetrate the inner regions. In other words, the ISM can self-shield. We have implemented this by exponentially suppressing the intensity of the UVB for gas that has $n_{\text{HI}} \geq 0.007 \text{ amu cm}^{-3}$ (De Rijcke *et al.*, 2013).

As such, the heating rate because of the UVB will be dependent on the gas density and on the time (or redshift), as well as the temperature and metallicity. Since the UVB will strongly influence the ionization state of the gas, the tables for the cooling rates, the mean constituent mass, and X_{ukin} will become 5-dimensional as well.

In Chapter 3 we did not include heating by the UVB, allowing us to only use the three-dimensional tables. Later on, we did include the UVB.

2.3.4 Star formation

Once a collapsing gas parcel becomes cold and dense enough, it is allowed to form stars. In other words, when

$$\nabla \cdot \mathbf{v} < 0 \quad (\text{Convergence criterium}), \quad (2.18)$$

$$T < T_{\text{crit}} = 15000 \text{ K} \quad (\text{Temperature criterium}), \quad (2.19)$$

$$\rho_{\text{g}} < \rho_{\text{crit}} = 100 \text{ amu cm}^{-3} \quad (\text{Density criterium}). \quad (2.20)$$

The density criterium (2.20) is the strictest one, as gaseous regions will be unlikely to reach the critical density if they are too warm or not collapsing. If a gas particle satisfies conditions (2.18)-(2.20), it is eligible to become a stellar particle, representing a single-age, single-metallicity stellar population. The actual star formation is implemented by a Schmidt law (Schmidt, 1959)

$$\frac{d\rho_{\star}}{dt} = -\frac{d\rho_{\text{g}}}{dt} = c_{\star} \frac{\rho_{\text{g}}}{t_{\text{g}}}, \quad (2.21)$$

with $\rho_{\star}$ and ρ_{g} respectively the stellar and gas density, $t_{\text{g}} = 1/\sqrt{4\pi G\rho_{\text{g}}}$ the dynamical time scale, and $c_{\star}$ the dimensionless star formation efficiency. We fix the latter at 0.25. Stinson *et al.* (2006) showed that varying this parameter between 0.05 and 1 had very little influence. This is because star formation is a self-regulating process.

2.3.5 Stellar feedback

The newly formed stellar particles are modelled with an initial-mass function (IMF), denoted here as $\Phi(m) = dN/dm$. It describes the probability that a star is born with a mass m in a stellar population. The Salpeter IMF (Salpeter, 1955)

2.3 Subgrid physics

$$\Phi(m)dm = Nm^{-2.35}dm, \quad (2.22)$$

is used in Chapter 3. The imposed lower and upper limits are respectively $m_l = 0.1 M_\odot$ and $m_u = 60 M_\odot$. $N = 0.06$ is a normalization factor so that $\int_{m_l}^{m_u} \Phi(m)dm = 1$.

In later chapters, we use the Chabrier IMF (Chabrier, 2003)

$$\Phi(m)dm = \begin{cases} N_1 m^{-1} e^{-A^2(\log_{10}(m)+B)^2} & \text{if } m \leq 1 M_\odot, \\ N_2 m^{-2.3} & \text{if } m \geq 1 M_\odot, \end{cases} \quad (2.23)$$

with $A^2 = 1.0502$ and $B = 1.1024$. $N_1 = N_2 e^{(AB)^2}$ so that both functions coincide for $m = 1 M_\odot$ and N_2 is again a normalization factor. For $m \geq 1 M_\odot$, the Chabrier IMF is roughly equivalent to the Salpeter IMF. The lower limit is the same as the one we used for the Salpeter IMF, but the imposed upper limit is now $m_u = 70 M_\odot$. We note that this gives rise to a small inconsistency in Chapter 3, since there we assumed a Salpeter IMF for the energy and metal injection, but the cooling rates assume a Chabrier IMF. However, the effects of this are small.

Young stellar populations give (thermal) feedback due to stellar winds (SW), supernovae type II (SNII), and type Ia (SNIa). The time intervals will differ for each type of feedback. For the SW, the interval will be from the birth of the stellar population until the last massive star turns SNII, which will have a mass $m = m_{\text{SNII},l} = 8 M_\odot$. The interval for SNII feedback is the time between the first and last SNII explosion. SNIa will happen in binary systems with a red giant (RG) and a white dwarf (WD) and is the result of the mass transfer from the RG on the WD when the former fills its Roche lobe. Only stars with masses between $m = m_{\text{SNIa},l} = 3 M_\odot$ and $m = m_{\text{SNIa},u} = 8 M_\odot$ are allowed to become SNIa. The main sequence lifetime of star is determined by its mass m . In Chapter 3, we use

$$\log t = 10 - 3.42 \log(m) + 0.88 \log^2(m), \quad (2.24)$$

as proposed by David *et al.* (1990), with t in years and m in $M_\odot$. For SNIa, we employ a delay time of 1.5 Gyr, which Yoshii *et al.* (1996) argue found to be the average needed accretion time for the WD to be pushed over its Chandrasekhar limit ($\sim 1.4 M_\odot$). This results in time intervals for

- **SW:** $0 - 4.3 \times 10^7$ yr
- **SNII:** $5 \times 10^6 - 4.3 \times 10^7$ yr
- **SNIa:** $1.54 \times 10^9 - 1.87 \times 10^9$ yr

From Chapter 4 on, we take a slightly different approach, since equation (2.24) behaves badly for high masses. The stellar main sequence lifetime is now calculated according to:

$$\log t = \begin{cases} 9.99 - 2.762 \log(m) & \text{if } m \leq 10 \text{ M}_\odot, \\ 7.99 - 0.764 \log(m) & \text{if } m \geq 10 \text{ M}_\odot. \end{cases} \quad (2.25)$$

Using this, a stellar population gives SNII feedback from 3.8×10^6 yrs after the birthtime of the stellar particle, until 3.1×10^7 yrs. SW again have the same upper limit as SNII and start from the birth of the stellar population. For SNIa, we now model the time interval as a Gaussian (Strolger *et al.*, 2010) centered around 4 Gyr and with a standard deviation of 0.8 Gyr (Bonaparte *et al.*, 2013).

SWs has an energy output of $E_{\text{SW}} = 10^{50}$ erg per massive star, while each SN event (of either type) has an energy output of $E_{\text{SN}} = 10^{51}$ erg (Thornton *et al.*, 1998). Our stellar particles represent a collection of stars, with masses determined by the assumed IMF. The total energy released by SNII from a stellar particle with mass M_{SSP} is thus

$$E_{\text{tot,SNII}} = E_{\text{SN}} \int_{m_{\text{SNII},l}}^{m_u} \Phi(m) dm \times \frac{M_{\text{SSP}}}{\int_{m_l}^{m_u} m \Phi(m) dm}. \quad = 0.191 \quad (2.26)$$

This energy is spread out uniformly over the time period where SNII feedback is given. The calculation of the total SW energy output is completely analogous. However, given that not every star in the $3 - 8 \text{ M}_\odot$ mass range will have a binary companion (that has become a RG by the time the primary star becomes a WD), we need to adjust the above formula with a factor A_{SNIa} to get the total SNIa energy output. We assume a ratio of 0.15 SNIa per SNII (Tsujiimoto *et al.*, 1995), which results in $A_{\text{SNIa}} = 0.0508$ for the recipes used in Chapter 3 and $A_{\text{SNIa}} = 0.0547$ for the later chapters.

The actual energy absorbed by the interstellar medium is determined by the feedback efficiency ϵ_{fb} . To produce sensible results, a high feedback efficiency is needed when using a high density threshold (Cloet-Osselaer *et al.*, 2012). We chose $\epsilon_{\text{fb}} = 0.7$. Along with the thermal energy from feedback, the gas particles are also enriched with metals produced in the stars and in supernovae.

As our simulations are not able to resolve the hot, low density cavities blown by supernovae, a gas particle that is receiving stellar feedback is not allowed to cool radiatively during that time step.

2.3.6 Mass return and chemical enrichment

After a star becomes a SN, it will leave behind a stellar remnant (i.e. black hole, neutron star or a white dwarf) with a lower mass than the original star. If we assume that the mass of a stellar remnant is constant, $M_{\text{rem}} = 1.4 M_{\odot}$, the mass fraction returned to the ISM by a SNII is thus

$$F_{\text{SNII}} = \frac{\int_{m_{\text{SNII},l}}^{m_{\text{SNII},u}} m\Phi(m)dm - M_{\text{rem}} \int_{m_{\text{SNII},l}}^{m_{\text{SNII},u}} \Phi(m)dm}{\int_{m_l}^{m_u} m\Phi(m)dm}. \quad (2.27)$$

Supernovae also pollute their surroundings with heavier elements. Knowing $M_i(m)$, the mass of an element i a star of mass m returns to the ISM in a SNII event (Nomoto *et al.*, 1997; François *et al.*, 2004), the total mass of element i injected in the ISM by a stellar population with mass M_{SSP} is given by

$$M_i = M_{\text{SSP}} \frac{\int_{m_{\text{SNII},l}}^{m_{\text{SNII},u}} M_i(m)\Phi(m)dm}{\int_{m_l}^{m_u} m\Phi(m)dm}. \quad (2.28)$$

For the returned mass fraction and metal yields by SNIa, one simply has to multiply (2.27) and (2.28) with A_{SNIa} , adjust the integral bounds in the numerators, set $M_{\text{rem}} = 0$ and now use the yields of Tsujimoto *et al.* (1995). Equation (2.28) simplifies for SNIa, since these metal yields are independent of progenitor mass.

2.3.7 Population III stars

Pop III stars formed in the early Universe out of pristine gas. To date no Pop III stars have been directly observed and their properties are necessarily derived from theory. Simulations predict Pop III star masses in the range between 0.7 and 300 $M_{\odot}$, with a significantly more top-heavy IMF than the Chabrier function of Pop II stars (Susa *et al.*, 2014, Figure 2.1). Compared with a Pop II particle, a Pop III star particle injects 4 times as much thermal energy into the ISM by SNII explosions and 40 times as much by young massive stars (Heger & Woosley, 2010). 45% of the mass of a Pop III particle is returned to the ISM, with the remainder locked up in remnants, while the chemical yields of Mg and Fe are taken to be 10% of the enrichment by normal SNII explosions (Heger & Woosley, 2010; Nomoto *et al.*, 2013). The latter is an approximation but, given the current theoretical uncertainties on Pop III yields and given the fact that a negligible fraction of the elements eventually present in the ISM comes from Pop III stars, this is without consequence. A star particle is treated as a Pop III particle if $[\text{Fe}/\text{H}] < -5$.

Recently, several other advanced treatment methods of rapid stellar feedback have been suggested and shown to be important in galaxy evolution models (e.g. Stinson *et al.*, 2013; Hopkins *et al.*, 2014). The main difference between these

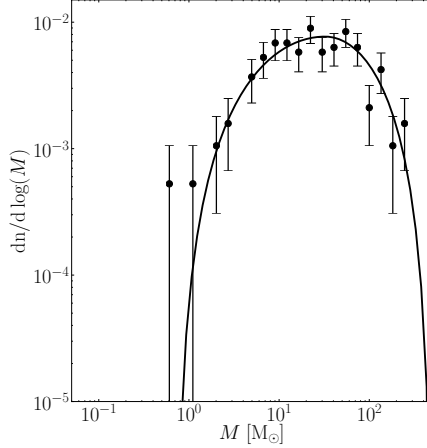


Figure 2.1: Initial mass function for the Pop III stars. We use a fit (solid line) through the data points (black dots) from Susa *et al.* (2014).

and the Pop III feedback proposed here is that the latter is inherently redshift dependent.

The motivation to include Pop III stars in our models is two-fold. We noticed that without it, the simulations form too many stars, based on stellar-to-halo-mass relations (e.g. Moster *et al.*, 2013), most of which formed at high redshift. Furthermore, the substructure of the ISM and the star formation rate at $z = 0$ (see Figs. 4.8 and 4.9, respectively) indicate that the feedback prescriptions for low redshifts are well motivated. We were thus specifically looking for a form of redshift-dependent feedback that reduces the amount of stars formed at high redshift and preserves the galaxy properties at low redshift.

Since I was personally strongly involved with the implementation of Pop III stars and the effects they have on the evolution of dwarf galaxies, they are more thoroughly discussed in Chapter 4.

2.4 Initial conditions

To construct the initial conditions for our simulations to start from, several approaches are possible.

2.4 Initial conditions

2.4.1 Cosmological simulations

The first approach is to start at high redshift with a large volume filled with a near-homogeneous matter distribution with over-densities determined by the fluctuations in the CMB (see e.g. Schaye *et al.*, 2015; Vogelsberger *et al.*, 2014). These are typically called *cosmological simulations* and have the most realistic initial conditions. The draw-back of this approach is that they are computationally very expensive and barely resolve fainter galaxies.

To alleviate the latter, *zoom simulations* can be set up by increasing the resolution of the particles within a certain smaller region of a cosmological simulation at $z = 0$ and resimulating this. This way, one can achieve a high resolution for a certain area of interest (e.g. a Milky Way analogue and its satellite system; Sawala *et al.*, 2016a) while maintaining the realistic initial conditions. However, zoom simulations remain computationally demanding. In this work, we do not run cosmological (zoom) simulations.

2.4.2 Isolated models

Another possible approach is to simulate a galaxy in isolation. The initial conditions for this are then typically a density profile with a certain prescribed form. These profiles are typically derived from cosmological simulations. One popular prescription for the dark matter halo density is the NFW profile (equation (1.5)). Next to dark matter, a galaxy also contains gas. The total gas mass is given by the the cosmic baryon fraction $f_{\text{bar}} = \frac{\Omega_{\text{bar}}}{\Omega_{\text{DM}}} \approx 0.17$ and will thus be fully determined by the mass of the dark matter halo. For the gas density profile, we could use a sphere with a uniform density or a the same profile as the dark matter density, scaled by f_{bar} . However, in this work, we will always use a pseudo-isothermal density profile (Revaz *et al.*, 2009):

$$\rho_g(r) = \frac{\rho_0}{1 + r^2/r_g^2}, \quad (2.29)$$

with ρ_0 the central gas density and r_g the scale length, determined in the same way as in Schroyen *et al.* (2013). Initially, there are no stars present in any of our simulations.

The gas sphere can be given a rotational velocity, according to

$$v_{\text{rot}}(r) = \frac{2}{\pi} \arctan\left(\frac{r}{r_{\text{rot,s}}}\right) v_{\text{max}}, \quad (2.30)$$

with v_{max} the initial rotational velocity at high radii and $r_{\text{rot,s}}$ the scale length of the rotational profile. This initial rotational velocity makes sure ordered motions are present in the gas from the start. Without this, the gas would be (mostly)

pressure supported. For all our simulations, we took $r_{\text{rot,s}} = 1$ kpc. Schroyen *et al.* (2011) compared simulations with rotation to those without and found them to be qualitatively very different. The most important finding was that in dwarf galaxies with significant rotation, the star formation is spread out almost homogeneously over the galactic body, as opposed to being localized to the centre. These initial conditions are idealized. However, using this approach, we have the advantage of being able to achieve a much higher resolution and a strong control on the properties of the simulated galaxy (most importantly, its dark matter mass).

GOGOIC² can be used to create the initial conditions, which is a wrapper program around KUZKUT³ and GANIC⁴ for the dark matter and gas halo respectively.

2.4.3 Merger trees

If we want to have more realistic initial conditions than the isolated models without having to rely on large cosmological simulations, we use the *merger tree* approach. This mass assembly history of a simulated dwarf galaxy is modeled as a merger tree constructed using the extended Press-Schechter formalism (Bond *et al.*, 1991; Parkinson *et al.*, 2008). We simulate the progenitors of the dwarf galaxy initially in isolation starting from $z = 13.5$ as described above but with a more generalized DM halo profile (Cen *et al.*, 2004):

$$\rho_{\text{DM}}(r) = \frac{\rho_0}{(r/r_s)^\alpha (1 + r/r_s)^{4-2\alpha}}, \quad (2.31)$$

with r_s the scale length, defined as the radius where the logarithmic slope of the density profile is -2 . The values for α and R_s are drawn from cosmologically motivated mass and redshift dependent probability distribution functions. We note that both the NFW profile as this one are *cuspy*, i.e. they are divergent for $r \rightarrow 0$. However, it was found that through gravitational interaction with the baryons, the dark matter halo will evolve from having a cusp to a core (Read & Gilmore, 2005; Cloet-Osselaer *et al.*, 2012; Teyssier *et al.*, 2013), in agreement with the observed cored dark matter profiles of dwarf galaxies (Salucci *et al.*, 2012). Hence, this density profile is representative only for the start of the simulations.

Subsequently, we let them coalesce as prescribed by the merger tree. The progenitors are given an initial rotation, with both the magnitude and direction of the initial angular momentum selected randomly. Each merger event is treated as a two-body interaction with orbital parameters drawn from probability distribution functions derived from cosmological simulations (Benson, 2005).

²<https://github.ugent.be/dwarf-galaxy-group/gogoIC>

³<https://github.ugent.be/dwarf-galaxy-group/kuzkut>

⁴<https://github.ugent.be/dwarf-galaxy-group/ganic>

2.5 Analysis

We do not explicitly include accretion in the simulations, however, we find that the simulated galaxies are surrounded by a very extended diffuse gas halo which does allow for smooth gas accretion implicitly in a self-consistent way. Similarly, the galaxies are embedded in an extended dark matter halo, which allows for growth of the dark matter halo through accretion.

2.5 Analysis

The post-processing analysis of the simulations was performed with our own publicly available software package `HYPLOT`⁵. This package has a GUI, which allows for a quick look at the snapshots of the simulations. However, `HYPLOT` can also be loaded in as a Python module, allowing one to make full use of its existing subroutines for calculating galaxy properties. Stellar luminosities are calculated using the `Miles` library, which gives the luminosity of a stellar population with a certain age and metallicity (Vazdekis *et al.*, 1996).

In Chapter 3, galaxy properties are generally obtained in a straight-forward way (e.g. their stellar mass is determined by summing the mass of each stellar particle). In the subsequent Chapters, we set out to determine the properties of the simulated galaxies in the same way as an observer would as closely as possible. This observationally motivated analysis is something that I made a significant contribution to and will prove to be of crucial importance throughout this work. The galaxy luminosities are now determined by fitting an exponential function to a simulated galaxy's surface brightness profiles out to R_{30} , the point where the surface brightness drops to $30 \text{ mag arcsec}^{-2}$, and extrapolating it to infinity to estimate the total luminosity. The V-band half-light radius R_e was determined from this fitted profile. $M_\star$ is estimated from the V and I-band luminosities (Bell & de Jong, 2001).

The neutral gas mass M_{HI} is determined by a straightforward summation of the masses of the gas particles multiplied with their neutral fraction. The UVB ionizes the outer, low-density regions of the ISM, limiting the H I to the more central, dense parts of the ISM of a simulated dwarf galaxy. It is important to note that the atomic hydrogen density of every gas particle has already been computed, based on its density, temperature, composition, and incident radiation field, to be used in the subgrid model of the simulations, as described in De Rijcke *et al.* (2013). Thus, all H I observables we describe below derive directly from the simulations without any extra assumptions or approximations.

The total baryonic mass is $M_{\text{bar}} = M_\star + M_{\text{gas,atomic}}$, with the atomic gas mass corrected for the Helium fraction: $M_{\text{gas,atomic}} = M_{\text{HI}}/(1 - Y_{\text{He}})$ (with the primordial Helium abundance $Y_{\text{He}} = 0.25$).

⁵<http://sourceforge.net/projects/hyplot/>

In order to produce an “observational” estimate for the circular velocity v_c , we fitted a gaussian to a mock H I spectrum of the galaxy viewed edge-on and adopted $v_c = W_{20}/2$, with W_{20} the full width at 20% of the maximum of this gaussian.

The cumulative star formation histories (CSFH) were determined from the birth times and masses of the stellar particles within a certain radius. These CSFHs, like the observed ones, thus do not take mass loss by SNe into account.

Using stellar evolution tracks for Pop II stars (Bertelli *et al.*, 2008, 2009) and for Pop III stars (Marigo *et al.*, 2001), the number of RGB stars within a given star particle can be calculated. Thus, the mean Iron abundance, $\langle[\text{Fe}/\text{H}]\rangle$, can be calculated via a sum over all stellar particles weighted by the number of RGB stars in each particle. This mimics the procedure actually followed for real dwarfs (Kirby *et al.*, 2013). For the gas, we computed the Oxygen abundance of dense gas, with each gas particle weighted by its ionized fraction. This is a measure for the metallicity of the ionized gas around star-forming regions. The solar logarithmic mass-fraction abundance of Iron is assumed to be $\text{Fe}/\text{H}_\odot = -2.756$ and the Magnesium to Iron ratio is $\text{Mg}/\text{Fe}_\odot = -0.261$ (Asplund *et al.*, 2009).

The Star formation rate (SFR) and atomic gas density, required for the Kennicutt-Schmidt relation, was determined within an aperture with a radius of 200 pc around the galaxy center. The 1D stellar velocity dispersion σ_* is measured along a line-of-sight through the galaxy, viewed edge-on, within an aperture of 1 half-light radii around the galaxy center.

The virial radius R_{halo} of a dark matter halo is the radius wherein the average density equals 200 times the mean density of the universe ($\rho_{\text{Univ}} = 9.47 \times 10^{-30} \text{ g cm}^{-3}$), while the virial mass M_{halo} is the total dark matter mass within R_{halo} : $M_{\text{halo}} = \frac{4\pi}{3} 200 \rho_{\text{Univ}} R_{\text{halo}}^3$.

In Chapter 5 further improved our observationally motivated analysis by making synthetic H I datacubes and using existing astronomical software to obtain resolved rotation curves for our simulated galaxies. These techniques are explained in detail in there Section 5.2.

Gaseous infall triggering starbursts in simulated dwarf galaxies

Sometimes science is more art than science, Morty. A lot of people don't get that.

Rick Sanchez

R. Verbeke, S. De Rijcke, M. Koleva, A. Cloet-Osselaer, B. Vandenbroucke and J. Schroyen (2014)

MNRAS, 420:1830

"Gaseous infall triggering starbursts in simulated dwarf galaxies"

3.1 Introduction

BLUE Compact Dwarfs (BCDs) are dwarf galaxies which are currently undergoing an intense period of star formation. The exact classification requirements for a dwarf galaxy to be a BCD differ from author to author. Zwicky (1964) first coined the term *compact*, in order to classify objects that could only just be distinguished from stars on photographic plates. Thuan & Martin (1981) defined a BCD as a galaxy with low luminosity ($M_B \geq -18$ mag), with small optical sizes (≤ 1 kpc), and an optical spectrum that exhibits sharp narrow emission lines superposed on a blue continuum, which are both generally caused by a high star formation rate (SFR). This is the reason BCDs were initially thought to be young objects going through their first episode of star formation. Deeper observations have shown that most BCDs in fact have a large population of old stars (Loose & Thuan, 1986; Kunth *et al.*, 1988; Papaderos *et al.*, 1996b; Cairós *et al.*, 2001). Loose & Thuan (1986) proposed a classification scheme based on the shape of the starburst region(s) and of the host galaxy.

- i0 BCDs show no evidence of having a host galaxy and are candidates of being true young objects.
- nE BCDs have a nucleated starburst component, embedded in a regular, elliptical host galaxy.
- iE BCDs also have an elliptical host galaxy, but the starburst component is irregular.
- iI BCDs have an irregular starburst embedded in an irregular host galaxy. This class can be further subdivided into
 - ▷ iI,C or cometary BCDs, which have very elongated host galaxies, with the star formation happening near one of the ends.
 - ▷ iI,M BCDs, which show clear signs of merging.

However, it is not clear if this classification has any physical meaning and a BCD is not necessarily clearly classified as one subtype.

Based on surface brightness profiles, Papaderos *et al.* (1996a); Salzer & Norton (1999) compared the structural parameters of BCD host galaxies to those of other dwarf galaxies and found that BCDs generally have smaller scale lengths and higher central surface brightnesses, meaning that their host galaxies are in general more compact than non-bursting dwarfs. Deeper observations have shown that the fainter, outer regions of BCDs (between 26 and 28 mag/arcsec²) do agree with dE/dIrr data (Micheva *et al.*, 2013a), although this might not hold for fainter BCDs (Micheva *et al.*, 2013b).

BCDs generally have a large central concentration of neutral gas (e.g. Taylor *et al.*, 1994, 1995; van Zee *et al.*, 1998; Simpson & Gottesman, 2000; Zhao *et al.*, 2013; Lelli *et al.*, 2014), although a variation in column density between different BCDs has been found (Simpson *et al.*, 2011). It was found that their gas can have very steep rotation curves, suggesting their dark matter is centrally concentrated as well (e.g. van Zee *et al.*, 2001; Lelli *et al.*, 2012a,b).

The gaseous metallicity of BCDs was found to generally be lower than that of dIs and dEs (e.g. Terlevich *et al.*, 1991; Izotov *et al.*, 1994; Izotov & Thuan, 1999; Hunter & Hoffman, 1999). Zhao *et al.* (2013) found that nE and iE BCDs are more metal-rich than iI BCDs, suggesting they might be at different evolutionary stages or have different progenitors.

Using an extended version of the GADGET-2 N-body/SPH code, we can self-consistently simulate isolated dwarf galaxies (Valcke *et al.*, 2008; Schroyen *et al.*, 2011; Cloet-Osselaer *et al.*, 2012; Schroyen *et al.*, 2013; Cloet-Osselaer *et al.*, 2014). In rotating models, the star formation is continuous while in the non-rotating dwarfs, we can find a SFR that goes up and down with a period of

3.2 Simulations

$\sim 10^8 - 10^9$ years (see also Stinson *et al.*, 2007; Revaz *et al.*, 2009). However, even in this breathing star formation, we can not identify true bursts.

Several external trigger mechanisms for the starburst in BCDs have been proposed and investigated. Tidal interaction could be the trigger (Brinks & Klein, 1988; Brinks, 1990; Campos-Aguilar & Moles, 1991; Noeske *et al.*, 2001), although not every BCD has a massive companion (Campos-Aguilar *et al.*, 1993; Telles & Terlevich, 1995). Mergers between gas-rich dwarf galaxies has also been suggested (Östlin *et al.*, 2001), which was confirmed as a plausible mechanism to trigger a starburst in simulated galaxies (Bekki, 2008; Di Matteo *et al.*, 2007; Cloet-Osselaer *et al.*, 2014).

As not all BCDs show signs of merging or tidal interaction, several internal mechanisms have been proposed as well, which are not present in our simulation code. Elmegreen *et al.* (2012) suggested in-spiraling clumps to drive dark matter, stars and gas from the halo towards the centre, which could feed the starburst. The stellar kinematics of BCDs can be very irregular, with the star forming regions standing out as dynamically distinct entities (Koleva *et al.*, 2014), suggesting they could be experiencing some form of interaction, even when galaxies seem to be in isolation. Starkenburg *et al.* (2016) found that the interaction between a gas-rich dwarf galaxy and a dark matter halo devoid of baryons can significantly increase the dwarfs star formation rate. External gas infall has been suggested as well, based on their high H I mass, disturbed and extended H I haloes and their low metallicities (Gordon & Gottesman, 1981; López-Sánchez *et al.*, 2012; Sánchez Almeida *et al.*, 2013; Zhao *et al.*, 2013; Sánchez Almeida *et al.*, 2014, 2015, 2016). H I clouds with no clear optical counterpart have been found around several BCDs (Taylor *et al.*, 1994; Hoffman *et al.*, 2003; Thuan *et al.*, 2004; Ramya *et al.*, 2009), so infall of such gas clouds is not an unlikely event.

In this Chapter, we shall use computer simulations to investigate whether a low metallicity gas cloud can trigger a starburst event in dwarf galaxies. This gas cloud can represent a truly external object or one that was formed within the galaxy’s halo.

An overview of the simulations we used is presented in Section 3.2. Section 3.3 gives an overview of our results, which are discussed in Section 3.4. We end with a summary in Section 3.5.

3.2 Simulations

3.2.1 The host galaxy

For the host galaxy, we use an isolated model, as laid out in Section 2.4.2. We start the simulation at redshift $z = 4.3$ and let it evolve for a period of 12.22 Gyr, until $z = 0$. We employ a flat, Λ -dominated cosmology with cold dark matter,

with $h = 0.71$, $\Omega_{\text{M}} = 0.2383$, $\Omega_{\text{DM}} = 0.1967$, $\Omega_{\text{baryon}}/\Omega_{\text{DM}} = 0.2115$.

The bulk of our dwarf galaxy models have a dark matter halo of $2.5 \times 10^9 M_{\odot}$ and an initial gas mass of $5.2 \times 10^8 M_{\odot}$ ("low mass dwarf galaxy models"). This will result in a galaxy with $-14 \text{ mag} \lesssim M_B \lesssim -12 \text{ mag}$ at redshift $z = 0$, where models with higher rotation are more luminous. To see how our results translate to higher luminosities, we also set up galaxies with dark matter mass $10^{10} M_{\odot}$ and initial gas mass $2 \times 10^9 M_{\odot}$ ("high mass dwarf galaxy models"), which will result in galaxies with $M_B \sim -15 \text{ mag}$.

We used 2×10^5 dark matter and 2×10^5 gas particles for the low mass models, resulting in dark matter and gas particles with masses of $\sim 12.5 \times 10^3 M_{\odot}$ and $\sim 2.6 \times 10^3 M_{\odot}$, respectively. 8×10^5 of each particle type were used for the high mass models, giving roughly the same mass resolution.

3.2.2 Gas clouds

To simulate a merger with a gas cloud, we first simulate a dwarf galaxy in isolation. Then, we introduce a gas cloud to the system at a certain time, a certain distance and with a certain velocity. We let it evolve until $z = 0$, allowing us to compare with observed BCDs at low redshift. While metallicity data exists for Milky Way HVCs, it is, at least to our knowledge, non-existent for gas clouds surrounding BCDs. Since BCDs are found predominantly in low-density environments such as the field, one would not expect gas clouds to have been significantly polluted. Therefore, the gas in the infalling gas cloud is initially given zero metallicity.

Mass

H I clouds with masses up to $\sim 10^7 M_{\odot}$ were found around BCDs (Hoffman *et al.*, 2003; Thuan *et al.*, 2004). For more estimates for the masses of infalling gas clouds, we could take a look at the masses of the high velocity clouds (HVCs) around the Milky Way and other giant spirals as more data is available for these (e.g. Oosterloo *et al.*, 2007; Lehner *et al.*, 2009; Putman *et al.*, 2009). The H I masses of these HVCs are estimated to range between $10^4 M_{\odot}$ and $10^7 M_{\odot}$, but with great uncertainties on these masses because of the uncertain distances (van Woerden & Wakker, 2004). A HVC with H I mass around $10^8 M_{\odot}$ has been found around M101, which most likely has a galactic origin (van der Hulst & Sancisi, 1988), so unlikely to be found around BCDs.

Kereš & Hernquist (2009) performed computer simulations of the formation of these HVCs around a Milky Way sized halo. These produced gas clouds with total gas masses (both neutral and ionized) ranging from $10^6 M_{\odot}$ to several $10^7 M_{\odot}$, with the lower limit being an artefact of the low resolution of the simulation.

3.2 Simulations

The clouds in this simulation located at high distances from the galactic centre have a trend to also be the more massive ones.

Therefore, although a mass of $\sim 10^7 M_\odot$ is, admittedly, at the high end of the gas cloud mass distribution (Pisano *et al.*, 2007), such massive clouds do exist. We expect that gas clouds with higher masses will have a bigger influence, so based on these observations and simulations, we will only discuss gas clouds with $M \approx 4 - 5 \times 10^7 M_\odot$.

The density of the gas cloud is chosen to have a $1/r$ -profile.

Radius

We investigated several radii for the gas cloud and separate the discussion qualitatively in compact gas clouds (CGCs) and extended gas clouds (EGCs). Following the Lemaître-Tolman-Bondi model for a spherical overdensity, we use for EGCs,

$$r_{\max} = \frac{0.09617}{5.55^{1/3}} \left[\frac{M_{\text{tot}}}{h^2(\Omega_m(1+z)^3 + 1 - \Omega_m)} \right]^{1/3} \text{ kpc}, \quad (3.1)$$

with M_{tot} the total mass of the gas cloud in $M_\odot$, $h = 0.71$, $z = 4.3$, and $\Omega_m = 0.2383$. E.g. for a gas cloud with mass $5.2 \times 10^7 M_\odot$, equation (3.1) gives us $r_{\max} \approx 7.5$ kpc.

The radius of a CGC we chose ad hoc, but significantly smaller than that obtained from equation (3.1). On the other hand, we also take it large enough to prevent that the density is too high and the gas cloud starts forming stars on its own. For CGCs we typically chose a radius of 2 – 3 kpc.

Temperature

As an extra condition, we want the gas cloud to remain stable over several Gyr, so that it does not start collapsing or expanding significantly. To do so, we set the temperature to be around the virial temperature

$$T_{\text{vir}} \simeq 6.8 \times 10^{-5} \mu \frac{M_{\text{tot}}}{r_{\max}} \frac{\text{kpc}}{M_\odot} \text{ K}, \quad (3.2)$$

with μ the mean molecular weight of the gas (for our purposes $\mu \approx 1.2$). We evolved isolated gas clouds during one Hubble time and found that the density profile within the inner ~ 5 kpc hardly changes. Only the tenuous outer layers expand because the gas clouds are set up with a vacuum boundary. For the actual simulations, we first let each gas cloud relax for 1 Gyr before putting them together with a dwarf galaxy in a merger simulation.

		EGC	CGC
Low mass	Slow rotation	B1	B2
	High rotation	B3	B4
High mass	Medium rotation	B5	B6

Table 3.1: Summary of the parameter space explored in the six models.

Velocity

The initial velocity given to the gas cloud was inspired by the escape velocity of the target galaxy, so the gas cloud seemingly falls in on radial orbits with an initial velocity

$$v_{\text{esc}}(R) = \sqrt{\frac{2GM}{R}}. \quad (3.3)$$

We put the centre of the gas cloud at $R = 30$ kpc, far enough from the galaxy. Looking into our dwarf galaxy simulations for the total mass enclosed with a radius of 30 kpc, we find $M_{<30\text{kpc}} \approx 2.5 \times 10^9 M_{\odot}$ for the low mass models and $M_{<30\text{kpc}} \approx 7 \times 10^9 M_{\odot}$ for the high mass models, which gives us $v_{\text{esc}} \approx 25$ km/s and $v_{\text{esc}} \approx 45$ km/s, respectively.

3.2.3 The models

Almost 150 simulations were run, all with variations in their set-up. The parameters that yield the biggest qualitative difference in the produced burst are the rotation rate of the host galaxy, the mass of the host and the radius of the gas cloud falling in. We will explicitly discuss 6 different simulations, in which these effects are very clear.

- B1 is a low-mass, slowly rotating host galaxy onto which an EGC falls in.
- B2 has the same host as B1, but a CGC is falling in.
- B3 is a low-mass host with fast rotation and an EGC falling in.
- B4 has the same host galaxy as B3, but with a CGC falling in.
- B5 is a high mass host and an EGC falling in.
- B6 has a high mass host galaxy, but with a CGC falling in.

3.3 Results

Name	$M_{\text{host,DM}}$ [$10^{10} M_{\odot}$]	$M_{\text{host,g}}$ [$10^{10} M_{\odot}$]	v_{max} [km s^{-1}]	M_{GC} [$10^6 M_{\odot}$]	GC type	R_{GC} [kpc]	t_{init} [Gyr]	(x, y, z) [kpc]	(v_x, v_y, v_z) [km s^{-1}]	SFR_{peak} [$10^{-2} M_{\odot} \text{ yr}^{-1}$]	b
B1	0.25	0.052	1.0	52	E	7.5	10.33	(-1,30,0)	(0,-25,0)	1.9	17.67
B2	0.25	0.052	1.0	42	C	2.0	10.33	(-1,30,0)	(0,-25,0)	1.1	9.73
B3	0.25	0.052	5.0	52	E	7.5	9.39	(-2,30,0)	(0,-25,0)	1.8	6.54
B4	0.25	0.052	5.0	42	C	2.0	9.54	(-2,30,0)	(0,-35,0)	3.0	10.71
B5	1	0.21	2.5	52	E	7.5	10.03	(-1,30,0)	(0,-45,0)	3.6	3.81
B6	1	0.21	2.5	42	C	3.0	10.13	(-2,30,0)	(0,-20,0)	4.9	5.16

Table 3.2: Columns indicate (1) model name, (2) dark matter mass of the host galaxy, (3) gas mass of the host galaxy, (4) the initial maximal rotational velocity of the gas (see eq. 2.30), (5) the mass of the gas cloud, (6) the type of gas cloud (E - extended or C - compact), (7) the radius of the gas cloud, (8) the time the gas cloud is introduced to the system, (9) the initial position of the gas cloud, (10) the initial velocity of the gas cloud, (11) the peak SFR for the merger simulation, (12) the burst factor (the peak in SFR divided by the SFR of the host galaxy in isolation, averaged over 3 Gyr).

Table 3.1 gives a short summary of this, while Table 3.2 gives a more detailed overview of the properties of the different models.

Other variations were explored, like the initial velocity of the gas cloud or the trajectory along which it falls in. These are not discussed explicitly using different models, but their results will be discussed where relevant in Section 3.3.

3.3 Results

3.3.1 Statistics

Before looking at the models presented in Section 3.2.3 and Table 3.2, we will discuss some general statistics of our simulations. For this, we define the burst-factor b as the peak SFR during the burst divided by the SFR of the isolated host averaged over the last 3 Gyr:

$$b = \text{SFR}_{\text{peak}} / \overline{\text{SFR}}_{\text{host}} \quad (3.4)$$

We will identify a burst if $b > 5$ and a strong burst if $b > 10$.

In total, 143 simulations were run. Out of these, 73 ($\sim 50\%$) had a burst, of which 9 were strong bursts. However, we can divide our simulations in several groups depending on the set-up, to distinguish several determining effects.

17 of the total simulation sample used a gas cloud with $M < 10^7 M_{\odot}$, of which none produced a burst.

In 8 of the remaining simulations, the gas cloud fell in along a prograde orbit, while in the others a retrograde orbit or a face-on collision was used. In only 1 of the former, a burst could be found. After close inspection of this simulation,

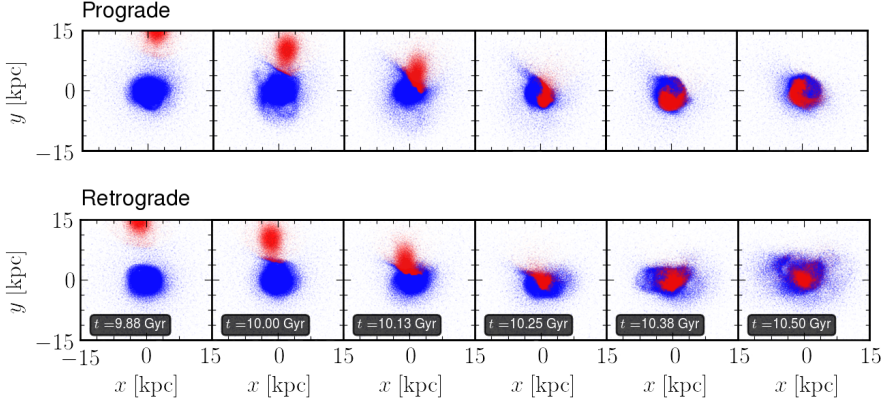


Figure 3.1: Gaseous infall on a dwarf galaxy. Red points show the gas cloud falling in, blue points represent the gas of the dwarf galaxy. The top panels show the set-up of B3: a gas cloud falling in on a dwarf galaxy along a prograde trajectory. The bottom panels shows the same set-up, but with the initial position of the gas cloud shifted to the left, resulting in a retrograde trajectory.

we found that the reason for the burst was that the gas cloud already had a high density on its own, out of which a large amount of stars could be formed when the gas cloud was moving through a deep gravitational potential and a high density region.

Of the remaining 118 simulations, 22 were high mass models and 96 were low mass models. 5 ($\sim 20\%$) of the high mass models had a burst (of which none were strong bursts), while 67 ($\sim 70\%$) of the low mass models had a burst (out of which 8 were strong bursts).

We do not wish to over-interpret these statistics, as they are not independent (e.g. several share the same host galaxy), but we can draw some general conclusions.

- We need a gas cloud with $M \gtrsim 10^7 M_\odot$ to trigger a burst in our models.
- Gas clouds falling in on a prograde orbit are unlikely to trigger a burst.
- For a gas cloud with the same mass, higher mass galaxies will be less likely to undergo a burst.

3.3.2 Gaseous infall and starburst modes

Gaseous infall can trigger enhanced star formation in a galaxy in several ways. This can be seen from Figure 3.2, where we show the star formation rate in

3.3 Results

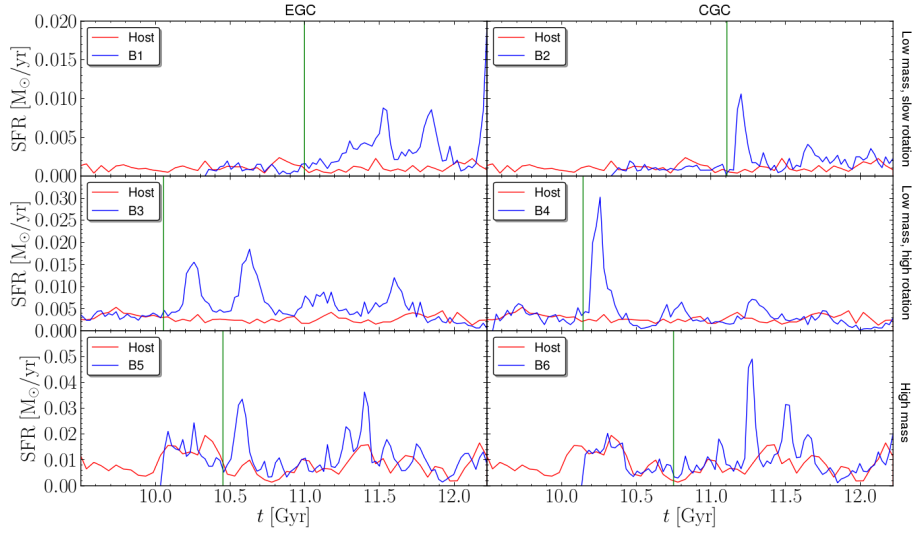


Figure 3.2: Star formation rate of bursting galaxies (blue), compared to that of their host galaxy in isolation (red). The models are the ones from Table 3.2. Green vertical lines roughly indicate when the gas cloud falls in on the dwarf galaxy.

function of time of our six models, compared to that of their host galaxies in isolation. The vertical lines roughly show the moments the gas cloud enters the system.

When a very extended gas cloud is falling in on a slowly or non-rotating host galaxy, it can compress the gas of the host out of which stars will be born (B5). However, this will not necessarily happen and we might not find a starburst initially. However, all the gas of the host galaxy will be affected, making it more turbulent and more likely to collapse, resulting in a strong burst at later times (B1).

In the case of a rotating galaxy (B3), the story is mainly the same, except that the direction the gas cloud is falling in from is important as it was concluded in Section 3.3.1. In a retrograde merger, the two gas reservoirs collide and lots of gas is driven towards the galaxy center. This is clearly in the left column of Figure 3.1. In a prograde merger, the velocities of the two gas reservoirs closely match and the infalling gas cloud swirls around the galaxy center. There is no significant angular momentum exchange between both gas reservoirs, no gas is driven towards the galaxy center, and no starburst is ignited. Movies of this can be found online².

For the high-mass models, these results remain valid, but since we use a gas cloud of the same mass as in the low mass models, the gas of the galaxy will be less disturbed in this case.

On the other hand, when we use a compact gas cloud, the gas cloud will only have an effect very locally. When stars are being formed as a direct consequence of the gas cloud impact, they will be formed along the path of the gas cloud. This is the case in B2 and B4. In the absence of rotation, the host galaxy will not be able to efficiently stop the gas cloud, so star forming region will have an elongated shape. Rotating galaxies will be able to stop it more efficiently, resulting in a more centralized star forming region.

In B6, the CGC did not immediately trigger an increased star formation rate. Rather, most of the gas flew through the galaxy, since the gas cloud has been significantly accelerated by the high mass dwarf galaxy. Shortly after however, the gas falls back, triggering the burst. The starburst is more centralized compared to B2 and B4 since the gas falling back in now has low densities. Thus the stars will be formed out of the old gas of the galaxy, much like the simulations with EGCs.

Figure 3.3 shows $B - I$ colour images of our models, illustrating these effects. These images were obtained by setting up Cartesian luminosity grids in both bands, to which we applied a Gaussian filter (with $\sigma = 100$ pc) to produce

²Youtube channel of our department: <https://www.youtube.com/user/AstroUGent>. A playlist with movies concerning this Chapter: http://www.youtube.com/playlist?list=PL-DZsb1G8F_kNcMwgkFNhQpHXRRhbOnCH5

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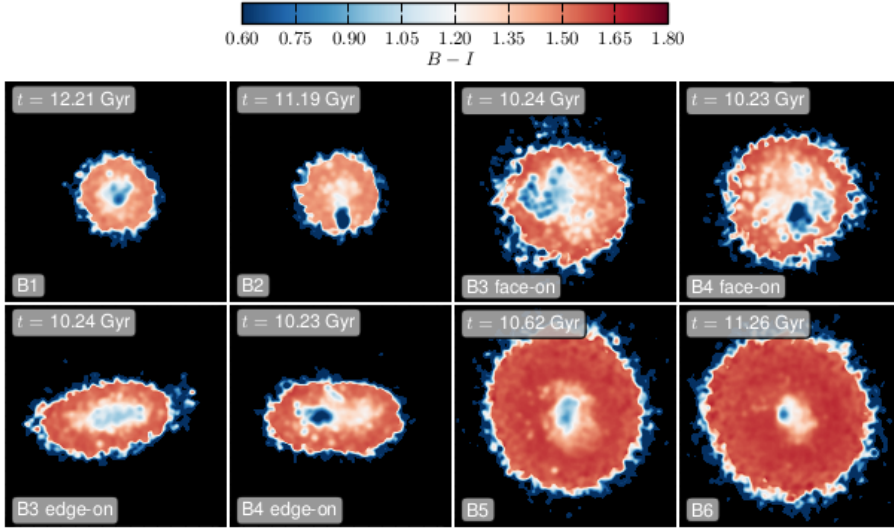


Figure 3.3: $B - I$ colour maps of our models during their bursts. All figures show 10 kpc side by side.

smoother images. The surface brightness of grid cells fainter than 29 mag/arcsec^2 were put to this magnitude. The blue edges are an artefact of this truncation, since the B-band will be rounded off more quickly than the I-band, so $B - I$ is underestimated.

3.3.3 Duration and periodicity of the starburst

After the initial burst, feedback in the form of SNII and stellar winds will efficiently remove the central gas reservoir and shut down the burst. Our bursts have lifetimes of a couple of 100 Myr, with a trend of stronger bursts having shorter lifetimes since the feedback will be more intense and remove the gas more efficiently. Indeed, in simulation B1 (top panel of Figure 3.2) we can identify a period of high star formation lasting $\sim 1 \text{ Gyr}$, shortly followed by a strong but short burst.

After the burst, we have little gas in the centre of the galaxy, making it easy for the gas to fall back in and produce secondary bursts. We find these subsequent bursts in simulations B1, B3, B5 and B6, as can be seen in Figure 3.2.

3.3.4 Strength of the burst

The mass of the gas cloud plays an important role in the strength of the burst. The more massive it is, the stronger the burst it can trigger. A gas cloud with $M = 5 \cdot 10^7 M_{\odot}$ can temporarily increase the SFR with a factor 20 in the low mass dwarf galaxy models.

For gas clouds of the same mass, the increase in SFR will be less significant in more massive dwarfs. These dwarf galaxies are already forming stars at higher rates because of the higher gas content and deeper gravitational potential, making it harder to get a relative increase in SFR similar to that in the less massive dwarf galaxy models. The more massive dwarfs will still have an increased SFR during the burst.

The peak in SFR of each model and the burstfactor b , which is the relative increase compared to the average SFR of the host over the last 3 Gyr, are given in column 11 and 12 in Table 3.2.

Secondary bursts can be stronger than the initial burst (e.g. B1) but generally, we can expect the strength of each subsequent burst to be lower than the previous one.

3.3.5 Metallicity

It is to be expected that the metallicity of the dwarf galaxy will drop when we let a metal-poor gas cloud merge with it. In all our models, we found that the metallicity of the gas within 5 kpc of the dwarf galaxy centre indeed drops significantly when the gas cloud enters the system. We show the evolution of the gaseous metallicity in the top panels of Figure 3.4 for B1 and B4.

However, looking at the B-band luminosity weighted metallicity of the stars (middle panel of Figure 3.4a), we see that the metallicity can increase significantly following the merger. In fact, the only ones of the models discussed here where the metallicity drops are model B2 and B4, as can be seen in Figure 3.4b for the latter.

In the other models, the gas falling in will not reach high enough densities to form stars at initial impact so if there is an immediate starburst (e.g. B5), the stars will be formed out of the old, metal enriched gas. In the case of bursts arising from the disturbed gas (e.g. B1), the gas will be well mixed and stars will be formed out of metal enriched gas as well. In B2 and B4, on the contrary, a significant part of the starburst is consuming the new, metal-poor gas.

Koleva *et al.* (2014) found that the starbursting regions can both be more metal-rich or metal-poor with respect to the host galaxy (their figures 3 and 4, respectively). Our results suggest that this might indicate whether the starburst is fueled from the old or the newly accreted gas.

3.3 Results

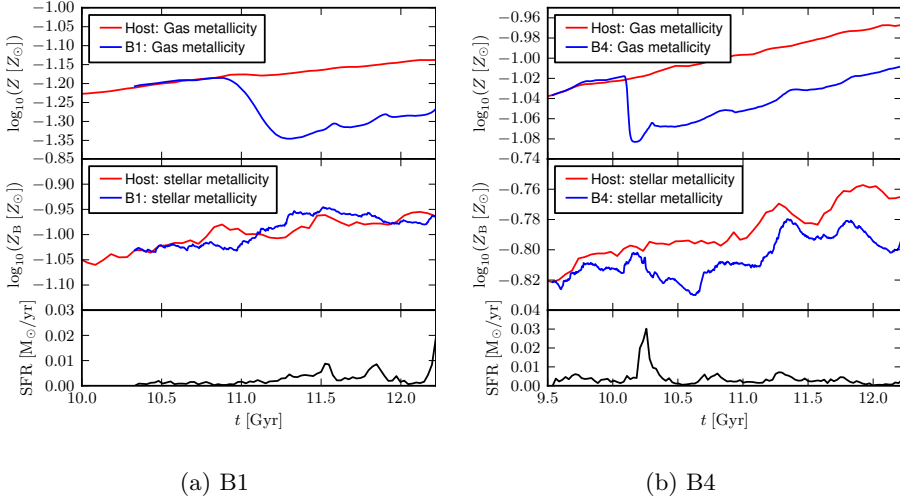


Figure 3.4: Evolution of the metallicity of the gas within 5 kpc (top panels), the B-band luminosity weighted metallicity of the stars (middle panels) and the SFR (bottom panels). Red shows the model in isolation, blue shows what happens when a gas cloud falls in.

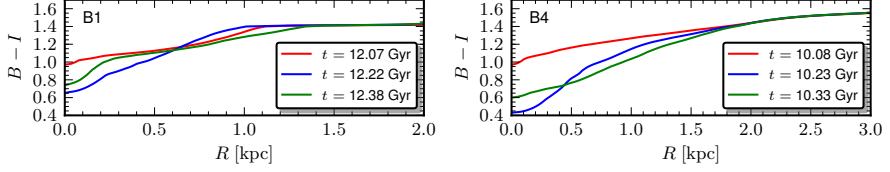


Figure 3.5: $B - I$ colour profile of model B1 (top) and B4 (bottom) before (red), during (blue) and after (green) their respective bursts.

3.3.6 Colour

Given the presence of irregular star formation knots in our galaxies during the starburst, we obtain colour profiles using a technique similar to *isophotal integration* for obtaining surface brightness profiles (Papaderos *et al.*, 1996b; Micheva *et al.*, 2013a,b). We start from the Cartesian grids used to produce the $B - I$ colourmaps in Figure 3.3. For each colour value, we determine the number of grid cells with colours bluer than this value and add their areas together. These areas are then converted to the equivalent radius of a circle with the same area. This method gives the radius in function of colour and it will produce a monotonically rising profile. The employed technique comes with the caveat that at some point, grid cells containing no stellar particles start influencing the profile. We cut off our profile before this happens. Also, the obtained radius can not be interpreted as a physical distance from the centre of the galaxy, unless in regular circular or elliptical systems.

The result for B1 and B4 can be seen in Figure 3.5, in which the colour profiles before, during and after their respective bursts are shown. We note that B1 has a strong burst right at the end of the simulation (top panel of Figure 3.2). To investigate what happens after the burst, we continued this simulation until $t = 12.7$ Gyr. For this simulation, we see a clear drop in colour for $R \lesssim 1$ kpc during the burst, with colour difference of over 0.3 mag in the centre, compared to before the burst. After the burst, the galaxy has become redder in the centre than during the burst, but is still bluer than before the burst. For $R \gtrsim 0.5$ kpc, the post-burst galaxy is bluer than during the burst. This can be due to recently formed stars moving to higher radii or stars being formed by the intense feedback following the central starburst. Since the stars are being formed from the metal-enriched gas of the dwarf galaxy, the bluer colours can be solely attributed by the age of the stellar population. For B4, the trend is the same, but we find much bluer colours: $B - I$ drops by ~ 0.6 mag in the centre, compared to before the burst. The reason is that now the stars are being formed from the metal-poor gas of the gas cloud.

3.3 Results

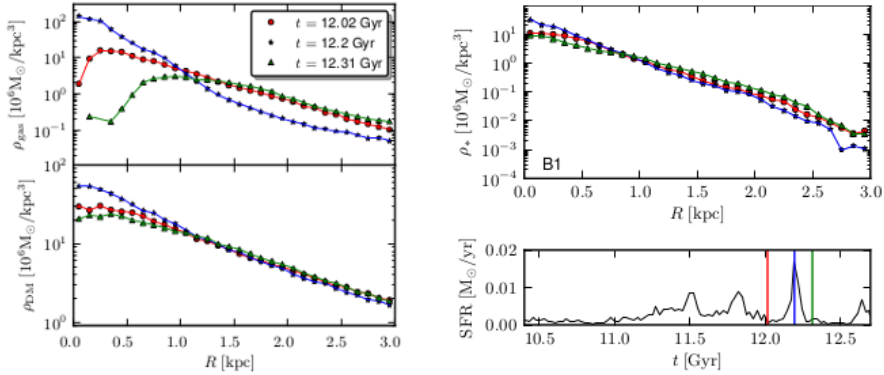


Figure 3.6: Gas, dark matter and stellar density profile of B1 before (red circles), during (blue stars) and after the starburst event (green triangles).

3.3.7 Density profiles

Figure 3.6 shows the density profile of the gas, dark matter and stars, at times before, during and after the burst. We see that for all the components, the central density rises significantly during the burst compared to before, while the density at higher radii drops. After the burst, the central density drops, even below the density before the burst, and gets higher in the outer regions. This effect is strongest for the gas.

No matter how the starburst is triggered exactly, the reason we have a burst is because the gas reaches a high density in a certain area. So the higher central gaseous densities should not come as a surprise. However, this large central gaseous mass causes the rest of the galaxy to contract. Hence we get a more compact galaxy, with a higher concentration of gas, dark matter and stars in the centre (also because of the high star formation) and a lower concentration in the outskirts of the galaxy. This redistribution of matter can in other words be attributed to the deepening of the gravitational potential. This process occurs in all models, albeit less significant in B2 and B4, where the impact of the CGC is a more local effect.

Shortly after the burst, the gas will receive a lot of feedback from the young stellar particles and will be blown to larger radii. The gravitational potential becomes shallower and the dark matter and stars also move outwards, resulting in a more diffuse galaxy. This process is present in all models and should be only dependent on the strength of the burst, as this determines the amount of feedback and thus the efficiency of the removal of the gas.

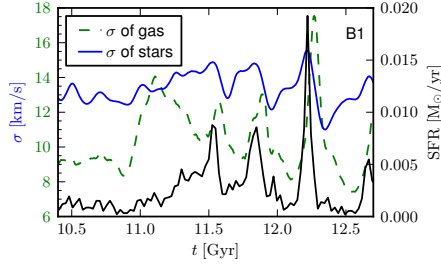


Figure 3.7: Velocity dispersion of the gas (green dashed line) and of the stars (blue solid line) compared to the SFR (black solid line) for simulation B1.

Since the expansion due to feedback will be faster than the previous collapse to the centre, the former will have a larger effect. So we are possibly left with a postburst galaxy that is more diffuse than preburst. This is very comparable to transforming a cuspy dark matter halo to a cored one through dynamical heating (Pontzen & Governato, 2012).

3.3.8 Velocity dispersion

We have a look at the velocity dispersion of both the gas (within 5 kpc) and the stars, defined as

$$\sigma = \sqrt{\frac{\sigma_x^2 + \sigma_y^2 + \sigma_z^2}{3}}. \quad (3.5)$$

In Figure 3.7, we show the evolution of σ of the gas and of the stars of B1, along with the SFR. When the gas enters the system (at ~ 11 Gyr), the velocity dispersion of the gas increases significantly. After it drops back, it follows the general trend of the SFR with a slight delay. The velocity dispersion of the stars increases and decreases together with the SFR. This trend is present in all our models, albeit less significant in the more massive ones B5 and B6.

Before the gas can form stars, it has to cool down and collapse, thus lowering its velocity dispersion. As new stars are being formed, gas with low σ is removed and, more importantly, the surrounding gas gets heated and it expands, increasing the gaseous velocity dispersion. This explains why the velocity dispersion of the gas follows the star formation rate with a slight delay.

On the other hand, as the gas collapses and produces a burst, the stellar body also becomes more compact and the velocity dispersion of the stars increases during a starburst.

Shortly after the burst, the stellar body is more diffuse and the velocity dispersion drops. This explains the concurrence of the stellar velocity dispersion curve and

3.3 Results

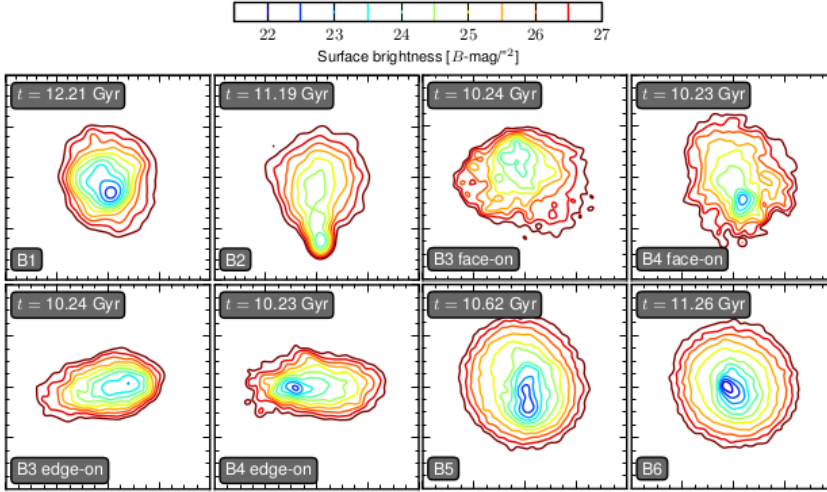


Figure 3.8: Isophotal maps in the B-band of our models during their burst. Sides for B1 and B2 are 4 kpc on both sides, for B3 and B4 6 kpc and for B5 and B6 both sides are 7 kpc. For B3 and B4, the models with a strongly rotating host galaxy, both the edge-on as the face-on maps are shown. Isophotes go from $\mu = 27$ mag/arcsec² (outermost and reddest) up to $\mu = 21.5$ mag/arcsec² (innermost and bluest), in increments of 0.5 mag/arcsec²

the SFR in Figure 3.7. The same is true for the dark matter. As an extra effect, we have a lot of young stars with low σ , further decreasing the velocity dispersion of the stars.

3.3.9 Isophotal maps

Figure 3.8 shows the isophotal maps of our models during their strongest peak in star formation. To obtain these, we produced a cartesian grid (with cells of width 70 pc), calculated the surface brightness in each cell, applied a Gaussian filter ($\sigma = 100$ pc) to smooth the data and plotted the contours. Isophotes go from $\mu = 27$ mag/arcsec² (most outer and reddest) up to $\mu = 21.5$ mag/arcsec² (most inner and bluest), in increments of 0.5 mag/arcsec²

Model B1 has a regular, slightly off-centre star forming region and the outer isophotes are nicely elliptical. Model B2 has a very off-centre star forming region, giving the entire galaxy a very elongated shape.

For B3 and B4, the models with a host with strong rotation, we show both a face-on and edge-on view. Face-on, both have very irregular outer isophotes,

which can be explained by the rotation (Schroyen *et al.*, 2011). The star forming region in both models is very off-centre. In B3, it is very irregular, while the star forming region in B4 is more regular. However, due to it being off-centre, it also gives the galaxy an elongated shape.

Edge-on, the outer isophotes are flattened and more regular. In B3, the star forming region is off-centre and slightly irregular, while in B4, it is more centralized.

Both the high mass models B5 and B6 show a regular, slightly off-centre star forming region embedded in an elliptical host galaxy, similar to UM452.

Comparing to the sample of Micheva *et al.* (2013b) (their figure 2), we find that B1 has the same characteristics as the UM491, B2 (and to a lesser extend, B3 and B4 viewed edge-on as well) resembles UM559, the face-on view of B3 and B4 looks qualitatively like UM461, and B5 and B6 resemble UM452.

3.4 Discussion

3.4.1 The triggering mechanism

Having established that an infalling gas cloud can trigger a significant increase in the star formation of a dwarf galaxy, we can wonder what exactly is the reason of this increase. In Figure 3.9, we show the radius and density of all gas particles in simulation B1 at different times. The gas already present in the dwarf galaxy is shown by yellow dots, while the blue dots represent the newly added gas of the gas cloud. The first panel ($t = 11$ Gyr) shows the extended gas cloud falling in. In the second panel ($t = 11.54$ Gyr), the influx of gas has led to the formation of many star-forming clouds and consequently to an increased SFR. Remarkably, virtually no “new” gas particles are mixed into these high-density, star-forming regions and the stars are being formed out of the “old” gas of the galaxy. This causes the stellar metallicity to further increase (Section 3.3.5). Clearly, this first sharp increase of the SFR is the direct result of the gas cloud compressing the old gas of the galaxy. This initial increase in SFR when the gas cloud enters the system, can be seen in all our models, except in B6. On the other hand, when the gas cloud is initially dense enough, stars can form out of the new gas, resulting in a more metal-poor starburst. This occurs in simulations B2 and B4.

Moreover, we can conclude from our simulations that gas cloud mergers also have a longer lasting impact on the SFR. The increased turbulence of the interstellar gas following a merger and starburst lasts for several Gyrs, as can be seen in the velocity dispersion of the gas (§3.3.8). During this time, the gas appears to be much more prone to gravitational collapse and therefore to subsequent starburst events. In the third panel ($t = 11.92$ Gyr) of Figure 3.9, the feedback following the previous star formation has blown out the gas from the central region, lowering

3.4 Discussion

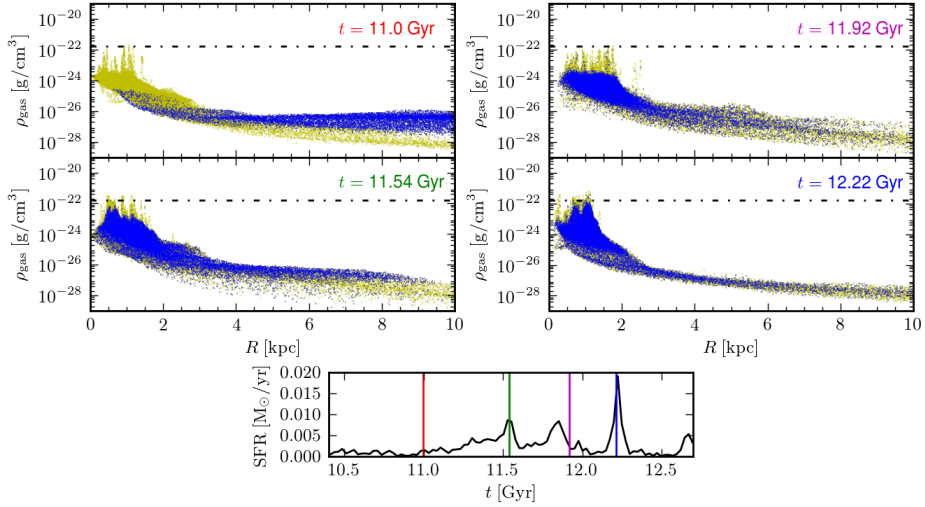


Figure 3.9: Scatter plot of the density and radius of all gas particles of simulation B1. Yellow points represent the old gas of the dwarf galaxy, blue points show the infalling gas cloud. The dash-dot line shows the density threshold for star formation.

the SFR. Around $R \sim 5$ kpc, for instance, one can see this gas being blown out. In the fourth panel ($t = 12.22$ Gyr), the gas has fallen back in and a starburst ensues, over a gigayear after the merger. Now, the gas is well mixed and the stars are being formed out of both the old and the newly added gas.

3.4.2 Surface brightness profile

As shown in Section 3.3.7, the stellar body of the galaxy becomes more compact during a starburst, in agreement with e.g. Papaderos *et al.* (1996b); Micheva *et al.* (2013a,b). They draw their conclusions by investigating the surface brightness profiles, usually obtained using *isophotal integration* due to the possible irregular shapes of BCDs. This method was already used to produce the colour profiles in Section 3.3.6, and now we obtain surface brightness profiles in a similar way. Figure 3.10 show the B-band surface brightness (SB) for model B1 and B6, before, during and after the starburst.

For the strong burst in B1 at $t \approx 12.2$ Gyr, we find an increase in SB in the centre, by about 1 mag/arcsec^2 , due to the bright, newly formed stellar populations. Moreover, as shown in §3.3.7, the outer stars are moving inwards causing the SB outside the inner ~ 0.5 kpc to decrease. This results in a overall steeper surface

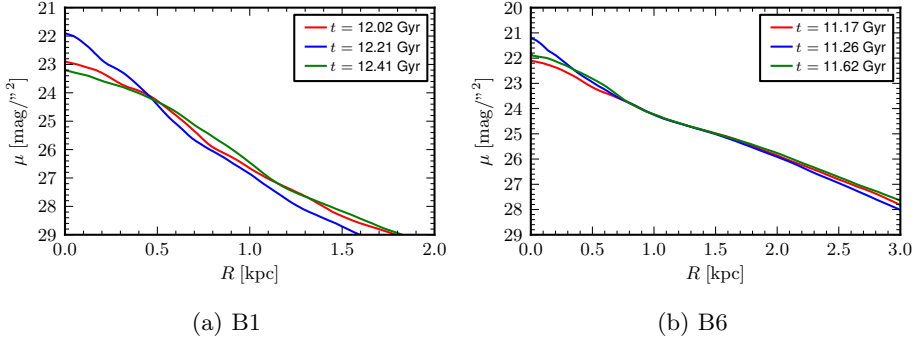


Figure 3.10: Isophotal surface brightness profile in the B-band before (red), during (blue) and after (green) the starburst.

brightness profile. For the more massive model B6, we also find an increased SB in the centre and a decreased SB at large radii, albeit less pronounced than in the low mass model B1. This is because the burst is weaker relative to the average SFR of the host galaxy (see Table 3.2).

3.4.3 Subtypes

Based on the isophotal maps discussed in Section 3.3.9, we can classify our models according to the classification scheme of Loose & Thuan (1986). Since our bursting models clearly have an old stellar population and we do not let two dwarf galaxies merge, we do not expect i0 or iLM BCDs to be found in our simulations. B1 and B6 have regular isophotes and a nuclear, albeit slightly off-centre, starburst and could thus be classified as nE BCDs. B5 has regular isophotes as well but the starburst is very off-centre and/or irregular, so we classify this model as a iE BCD. When viewing B3 and B4 face-on, we see an irregular, off-centre starburst in a highly irregular host. One would classify these as iI BCDs. However, when B3 and B4 are viewed edge-on, the starburst region can be significantly off-centre, giving these models a flattened, cometary look. For this viewing geometry, B3 and B4 would be classified as a iLC. The same classification applies to B2.

Based on these findings and the qualitative differences between the models, we can propose an explanation and make predictions for the different subtypes and their properties:

- iE BCDs can be produced when the infalling gas cloud immediately triggers a burst in a not, or only slowly, rotating dwarf, as the dense gas will most likely

3.4 Discussion

not be centralized and have irregular shapes and the outer isophotes will have regular shapes (e.g. models B2, B5). Of course, if the star-forming region is significantly off-centre, the galaxy might be labelled as a cometary BCD.

- nE BCDs on the other hand will more likely be found in secondary bursts in the same host galaxy (e.g. models B1, B6) as gas has re-accumulated at the galaxy centre and ignited in a new starburst. However, the supernova feedback from this central starburst can lead to star formation in dense clouds on the edges of supernova blown bubbles, resulting in more irregular star forming regions. Thus, a nE BCD may evolve into an iE BCD.
- ii BCDs are found in rotating host galaxies (e.g. models B3, B4). Schroyen *et al.* (2011) found that in simulated dwarf galaxies with a significant amount of rotation, star formation occurs continuously over the galactic body, in contrast to a more bursty, centralized star formation in dwarfs with slow (or no) rotation. This more dispersed star formation results in an irregular stellar body. So when the galaxy enters the starburst phase, the star forming region will most likely be off-centre and have an irregular shape. When looking at an ii BCD with off-centre star formation, it could be classified as an iI,C BCD, when viewed edge-on.

Our other simulations which produced a burst but which were not explicitly discussed here, all agree with this explanation for the classification of the different subtypes as well.

There are indeed a number of observations that support this picture. ii BCDs were found to be in general bluer (Gil de Paz *et al.*, 2003), more gas-rich, and more metal-poor (Zhao *et al.*, 2013) than nE/iE BCDs and their H I distribution is more clumpy rather than centralized (Taylor *et al.*, 1994; van Zee *et al.*, 1998; Thuan *et al.*, 2004). All of these findings can be explained by the degree of rotation they have (Schroyen *et al.*, 2011). Thuan *et al.* (2004) found that the nE and iE BCDs in their sample show no regular rotation, while the two cometary BCDs in their sample are rotationally dominated. At least 5 of the 7 iI,C BCDs investigated in Sánchez Almeida *et al.* (2013) are rotating. Observations have shown that some iE BCDs have their star forming regions organized in a ring (Gil de Paz *et al.*, 2003), supporting the idea of propagating star formation and an evolution from nE to iE BCD.

Noeske *et al.* (2000) found that the host galaxies of iI,C BCDs generally have structural parameters that lie between those of BCDs and other dwarfs. If our models are indicative for real life BCDs, the fact that these are less compact than nE/iE BCDs can be explained in two ways. Firstly, if star formation occurs very off-centre, there will not be such a large increase in the central mass and thus the gravitational potential will deepen less. This results in a less compact host galaxy. Secondly, iI,C BCDs are likely to be found in rotating host galaxies,

which are generally already more diffuse. On the other hand, when looking at a flattened galaxy edge-on, the surface brightness will drop more rapidly, resulting in a seemingly more compact galaxy, as opposed to when we would view it face-on. We thus propose that *ii* BCDs should be even more diffuse than *iI,C* BCDs, and their hosts will have scale lengths and central surface brightnesses in the parameter space occupied by *dE* and *dIs*. To the best of our knowledge, no large sample study has been performed of the structural parameters of the host of *ii* BCDs. The sample of Micheva *et al.* (2013a,b) contains some *ii* BCDs, who do indeed tend to lie in the parameter region of the *dEs* and *dIs*.

3.4.4 Gas kinematics

As the gas and stellar component of the galaxy becomes more centralized, we expect that their rotational velocity will increase due to conservation of angular momentum. This would explain the observations by van Zee *et al.* (2001), who found that the *H I* of BCDs have steeply rising rotation curves. Figure 3.11 shows the 2-dimensional gaseous line-of-sight velocity fields of B1 and B4 (edge-on view), left and right respectively, during the burst, overlaid with the *B*-luminosity contours. We show these two quantities along a slit placed along the major axis to mimic how this would be detected observationally. For B1, we indeed find a steeply rising rotation curve.

For B4, we see that the galaxy overall has solid body rotation around the minor axis (the *z*-axis), but the kinematics of the starburst region are very different (as found by e.g. Östlin *et al.*, 2004; Koleva *et al.*, 2014). In this model, a CGC has fallen in, coming from the right side in Figure 3.11. This impact will cause a significant amount of gas in the red region (with positive v_x) to move to the blue region (with negative v_x), causing a disturbance in the solid body rotation pattern of the galaxy. Looking along the *y*-axis (along which the gas cloud fell in), we will see a similar disturbance while in the *z*-axis, no significant difference will be found.

We can conclude that if the burst happens right after the gas cloud fell in and depending on our viewing angle at the galaxy, the gas can have very disturbed kinematics. In secondary bursts on the other hand, like in B1, the gas will most likely show no disturbance.

3.4.5 Stellar kinematics

Koleva *et al.* (2014) found that the star forming regions of BCDs can have kinematics that are very different from those of their host galaxies. The top panel of Figure 3.12 shows the velocity fields of the stars of simulation B2. The radial velocity and surface brightness along a slit along the major axis is shown in the bottom panel. As we can see, the velocity of the gas cloud that fell in results in

3.4 Discussion

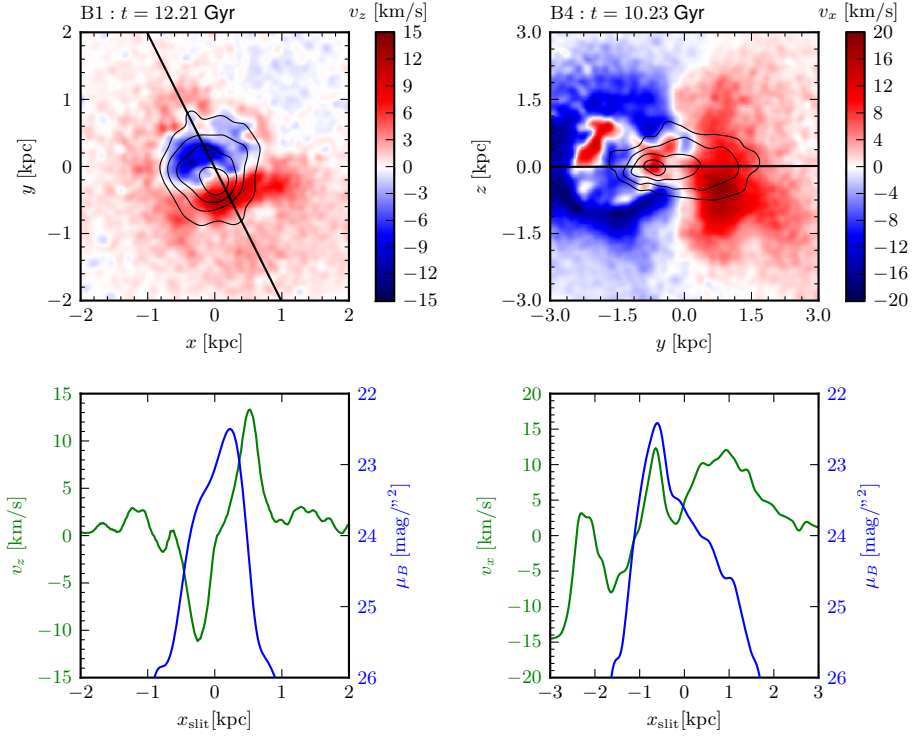


Figure 3.11: Top: velocity field of the gas overlaid with the surface brightness contours in the B-band (top). Bottom: the line-of-sight velocity (green) and the surface brightness (blue) along a slit put on the major axis for simulations B1 (left) and B4 (right).

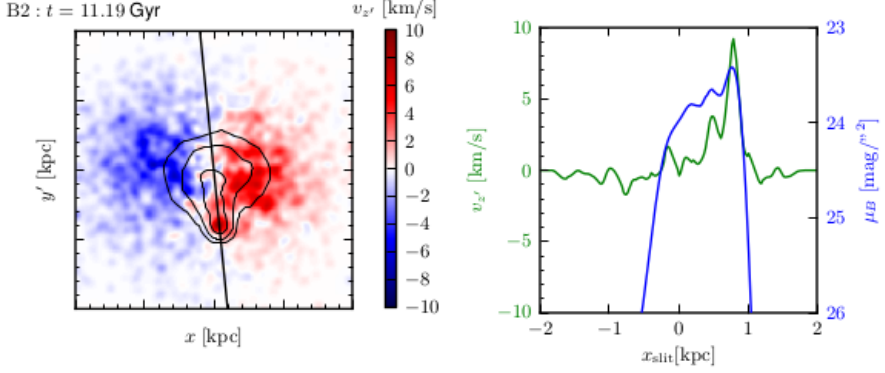


Figure 3.12: Same as in Figure 3.11 but for the stars of simulation B2. Note that the simulation was rotated $\sim 50^\circ$ around the x -axis to show the effects more clearly.

a star forming region with different kinematical properties than that of the host galaxy. Thus, our simulations are able to reproduce the findings of Koleva *et al.* (2014), at least for simulations with a compact infalling gas cloud, as in B2.

3.4.6 Evolutionary path

It has been long discussed what BCDs will look like when their starburst has faded. Among others, Papaderos *et al.* (1996b) suggested that postburst dwarf galaxies can not reside in the general dE or dI population without a drastic change in their structural parameters. Indeed, if the stellar body is assumed to be static, fading of the stellar populations is too slow to transform the typical surface brightnesses of BCDs to those typical for dIs. van Zee *et al.* (2001) state that a passive evolution to dE is unlikely as it would require a significant loss of angular momentum. They suggest that so-called compact dI (e.g. van Zee, 2000; Patterson & Thuan, 1996) are candidate faded BCDs, since they have similar scale lengths and have significant colour gradients.

Looking at our models, we find that the stellar body is not static and generally postburst galaxies tend to be more diffuse after the burst than during the burst, when the galaxy would be classified as a BCD. Often, the galaxy is even more diffuse than before the burst(s), placing it within the structural parameter space occupied by dEs and dIs. Moreover, as the galaxy re-expands after the burst, the steep rise of its rotation curve disappears (as can be seen in Figure 3.13a),

3.4 Discussion

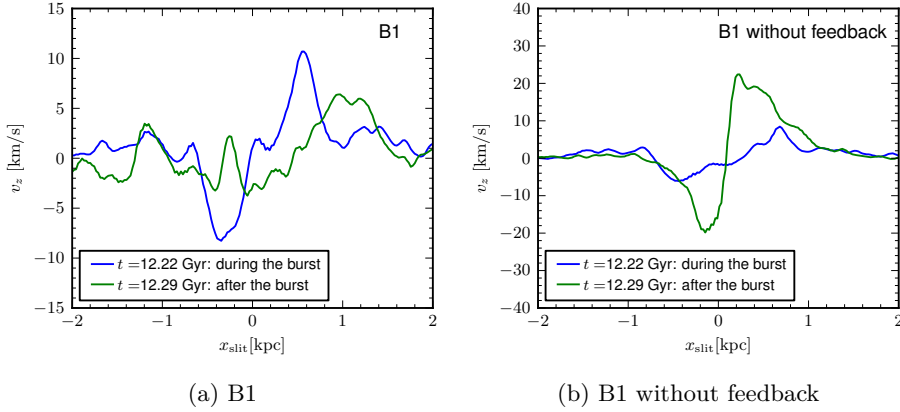


Figure 3.13: **Left:** line of sight velocity of the gas along the same slit as in Figure 3.11 for simulation B1 during (blue line) and after (green line) the starburst. **Right:** same, but for a model with both star formation and feedback turned off after the burst. Note that the vertical scale is different in the left and right panels.

removing also this obstacle to a possible BCD-dI transition. On top of that, the time-scale of a burst is much smaller than the gas exhaustion time and the feedback following the starburst is not strong enough to completely remove all the gas of the galaxy. Thus BCDs that were dIs before the burst will be dIs again after the burst.

3.4.7 The role of feedback

Stellar feedback through supernovae and stellar winds is assumed to play a vital role in quenching the starburst. Furthermore, it seems that the strong feedback effects result in a diffuse post-burst galaxy, with shallow rotation curves and low stellar velocity dispersion, as discussed in Sects. 3.3.7, 3.4.6 and 3.3.8. To confirm that this evolution is primarily due to feedback, we have rerun simulation B1, but with feedback turned off after $t = 12.22$ Gyr (the time of the burst). Since there would be no way of quenching the starburst, we also turn off star formation.

There is now nothing to prevent the gas from collapsing further, resulting in a very deep gravitational potential, a very compact galaxy (in both gas, stars and dark matter), high velocity dispersion and a very steep rotation curve. The latter is shown in Figure 3.13b.

3.5 Summary

We have run a large suite of numerical simulations of gas clouds falling in on dwarf galaxies, using different galaxy and cloud masses and sizes, and different orbital parameters. Our goal was to test whether starbursts could be ignited this way and whether the resulting galaxies could explain some of the properties of observed Blue Compact Dwarfs.

Based on our simulations, we conclude that infalling gas can indeed trigger a starburst in dwarf galaxies by significantly disturbing the gas, especially in dwarf galaxies with $M_B \gtrsim -14$ mag. We found that gas cloud masses significantly below $\sim 10^7 M_\odot$ do not trigger starbursts in the mass regime of the dwarfs simulated by us. If such clouds are rare, then this tells us that, although this is a viable way of triggering starbursts in dwarfs, this is not the triggering mechanism in the majority of the BCDs. Of course, less massive clouds could still trigger starbursts in dwarf galaxies with lower masses than the ones we simulated. The starburst phase seems to be episodic, alternated with periods of little star formation. Following the infall of metal-poor gas, the overall gas metallicity of the galaxy drops. However, the stellar metallicity can either increase or decrease, depending on the source of the gas from which the stars are formed: the pre-enriched galaxy or the metal-poor infalling gas cloud. In most of our models, the stars were formed from the old gas, resulting in metal-rich stars. Only in models with compact infalling gas clouds can stars form from the newly arrived gas, leading to more metal-poor newborn stars. Following the impact of the gas cloud, the gas kinematics can become very disturbed. Stars that form from kinematically distinct gas clumps tend to be also kinematically peculiar when compared with the stellar body of the galaxy.

As gas flows towards the galaxy centre to fuel the central starburst, the gravitational potential deepens. In response, the whole body of the galaxy becomes more compact and the rotation curve steepens. Once gas is rapidly evacuated from the galaxy centre by supernova feedback, the galaxy re-expands and becomes more diffuse again, flattening the rotation curve. This removes some of the arguments against evolutionary links between BCDs and dIIs.

As discussed in paragraph 3.4.3, we can attempt an explanation for the physical differences between the subtypes of BCDs in the classification scheme of Loose & Thuan (1986) within the scope of our simulations:

- nE and iE BCDs have host galaxies that are slowly (or not) rotating. nE BCDs can evolve into iE types as star formation propagates through the host galaxy.
- ii BCDs have flattened host galaxies with strong rotation, viewed approximately face-on.

3.5 Summary

- iI,C BCDs have flattened, strongly rotating host galaxy, viewed approximately edge-on.

How the First Stars Shaped the Faintest Gas-Dominated Dwarf Galaxies

Let them come. There is one
Dwarf yet in Moria who still
draws breath.

Gimli, son of Glóin

R. Verbeke , S. De Rijcke and B. Vandenbroucke (2015)

ApJ, 815:85

"How the first stars shaped the faintest gas-dominated dwarf galaxies"

4.1 Introduction

LEO P (Giovanelli *et al.*, 2013) and Pisces A (Tollerud *et al.*, 2015) are the latest additions to a growing list of faint, gas-rich, isolated dwarf galaxies; a list that dates back at least to the discovery of Leo T (Irwin *et al.*, 2007). These galaxies have luminosities of only a few $10^5 L_{\odot}$ (Ryan-Weber *et al.*, 2008; McQuinn *et al.*, 2013; Tollerud *et al.*, 2015) and H I masses 3 – 4 times higher than their stellar mass. Rotation velocities, derived from radio observations of the H I, are estimated at $v_{\text{rot}} \sim 12.5 \text{ km s}^{-1}$ (Leo T: Ryan-Weber *et al.*, 2008), $v_{\text{rot}} \sim 17 \text{ km s}^{-1}$ (Pisces A: Tollerud *et al.*, 2015) and $v_{\text{rot}} \sim 18 \text{ km s}^{-1}$ (Leo P: Giovanelli *et al.*, 2013; Bernstein-Cooper *et al.*, 2014). The available data suggest that all three dwarfs formed stars continuously, although at a very low and highly variable rate (Clementini *et al.*, 2012; Weisz *et al.*, 2012; McQuinn *et al.*, 2013, 2015).

The existence of these systems poses a strong challenge to Λ CDM, the standard model for galaxy formation and evolution: cosmological simulations predict that

half of the galaxies in the nearby Universe with a circular velocity $\sim 25 \text{ km s}^{-1}$ are dark; these halos were never able to form stars. At a circular velocity of $\sim 15 \text{ km s}^{-1}$, comparable to Leo T, Leo P, and Pisces A, over 90% of all halos is predicted to be dark (Sawala *et al.*, 2016b). Indeed, supernova explosions together with the cosmic ultraviolet background (UVB), produced by the first galaxies and quasars (Efstathiou, 1992; Haardt & Madau, 1996; Faucher-Giguère *et al.*, 2009), tend to remove the star-forming gas from low-mass dwarf galaxies. However, Tollerud *et al.* (2015) found that an order of magnitude difference between the observed number density of H I-detected faint dwarfs and that of corresponding dark-matter halos predicted from cosmological simulations is highly unlikely.

Clearly, these are not rare objects and many more such faint systems are likely awaiting discovery (Adams *et al.*, 2013; Faerman *et al.*, 2013; Adams *et al.*, 2015; Janesh *et al.*, 2015; Sand *et al.*, 2015). Thus, there appears to be a disagreement between the predicted and the observed abundance of faint, gas-dominated, star-forming dwarf galaxies near the Local Group.

The process of galaxy formation is very challenging to model and the observable properties of simulated galaxies will strongly depend on the chosen set of parameters. While parameters may be tuned to get one or several galaxy properties in agreement with observations, reproducing a broad range of them is not a trivial task. Failing to reproduce one or more observable properties may be indicative that an important astrophysical process is not taken into account in the models. We thus argue that to truly reproduce realistic galaxies, one has to look at all the known observable galaxy properties. Many efforts have been made to self-consistently simulate the formation and evolution of low-mass dwarf galaxies (e.g. Governato *et al.*, 2010; Munshi *et al.*, 2013; Cloet-Osselaer *et al.*, 2014; Shen *et al.*, 2014; Benítez-Llambay *et al.*, 2015; Oñorbe *et al.*, 2015; Sawala *et al.*, 2015; Trujillo-Gomez *et al.*, 2015; Wang *et al.*, 2015). In this and the following Chapter, we compare our simulations with as many observable properties as possible.

Here, we propose a theoretically attractive way of alleviating these tensions between simulations and observations: the energetic feedback from the first stars that formed in the universe. These so-called “population III”, or Pop III, stars are expected to have very different properties than stars born out of even very weakly enriched gas, which are called “population II”, or Pop II, stars. As we will show, this allows faint dwarfs with stellar masses of $M_{\star} \sim 10^{5-6} M_{\odot}$ to grow in dark-matter halos of $M_{\text{DM}} \sim 10^{9-10} M_{\odot}$ which are massive enough to retain their cold gas reservoirs.

In Section 4.2, we give an overview of the simulations we used. The results of our simulations are shown and discussed in Section 4.3. Section 4.4 provides a short summary of this work.

4.2 Simulations

(1) Name	(2) $M_{\text{DM,tot}}$ [$10^9 M_\odot$]	(3) M_{prog} [$10^9 M_\odot$]	(4) M_{res} [$10^7 M_\odot$]	(5) n_{DM} $\times 10^5$	(6) n_{SPH} $\times 10^5$	(7) ϵ_f [pc]	(8) Symbol	(9) Notes
DG9e9	9	1.36	5	4.5	18	6.15	★	
DG10e9	10	1.82	5	5	20	6.15	○	
DG12e9	12	1.88	5	6	18	6.77	◁	
DG13e9a	13	1.86	5	6.5	6.5	9.76	▽	
DG13e9b	13	2.04	5	6.5	6.5	9.76	△	Different initial conditions
DG15e9a	15	1.82	5	7.5	7.5	9.76	□	
DG15e9b	15	1.57	5	7.5	7.5	9.76	◇	Different initial conditions
DG20e9	20	3.07	10	10	10	9.76	○	
DG50e9	50	2.98	10	10	10	13.25	○	
DG1e11	100	7.34	20	20	20	13.25	○	
DG15e9b-CT	15	1.57	5	7.5	15	7.75	◇	Convergence test of DG15e9b
DG10e9-NP3	10	1.82	5	5	5	9.76	○	DG10e9 without Pop III feedback
DG12e9-NP3	12	1.88	5	6	6	9.76	◁	DG12e9 without Pop III feedback
DG13e9b-NP3	15	2.04	5	7.5	7.5	9.76	△	DG13e9b without Pop III feedback

Table 4.1: Overview of the simulations discussed in detail in this chapter. (1): the name of the simulation, (2): total dark matter mass in the simulation, (3): the dark matter mass of the most massive progenitor, (4): the minimal progenitors mass in the merger tree, (5): number of dark matter particles, (6): number of gas particles, (7): force resolution, (8): symbol used in the figures, (9): additional notes.

4.2 Simulations

4.2.1 Overview

We use the MORIA (Models of Realistic dwarfs In Action) simulations to test our hypothesis. This is a suite of (currently) ~ 30 simulations covering the dwarf mass regime ($M_\star \sim 10^{5-9} M_\odot$) with different mass assembly histories in a cosmological setting with added Pop III feedback. This allows for a comparison with some of the faintest dwarf irregulars, like Leo T, Leo P, and Pisces A. 10 of the MORIA simulations are discussed in more detail. We also present convergence test simulations and control simulations with the same initial conditions and mass assembly histories but without Pop III feedback. We refer to Table 4.1 for an overview of the 10 selected simulations. Figure 4.1 shows the different assembly histories of the selected simulations.

4.2.2 Resolution and convergence test

All but two simulations were run with the mass of the dark matter particles fixed at $m_{\text{DM}} = 2 \times 10^4 M_\odot$, except for DG50e9 and DG1e11 with $m_{\text{DM}} = 5 \times 10^4 M_\odot$. In most simulations, the mass of the gas particles was simply scaled according to the assumed baryonic to dark-matter density ratio: $\Omega_{\text{bar}}/\Omega_{\text{DM}} = 0.2115$. Exceptions to this are DG10e9, DG12e9 and DG15e9b-CT, in which the gas resolution was increased respectively with a factor of 4, 3 and 2. In the latter,

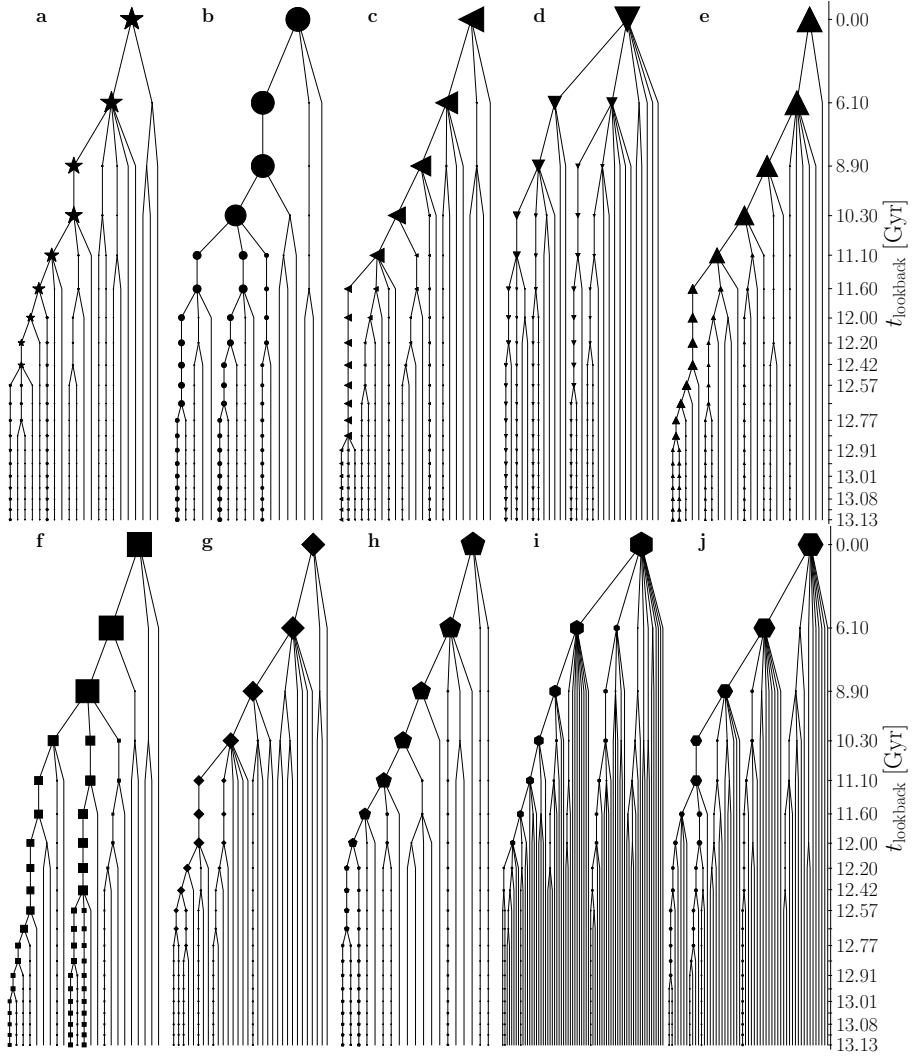


Figure 4.1: Merger trees of the simulations. (a) DG9e9, (b) DG10e9, (c) DG12e9, (d) DG13e9L, (e) DG13e9E, (f) DG15e9L, (g) DG15e9E, (h) DG20e9, (i) DG50e9 and (j) DG1e11, and their respective re-simulations without Pop III stars and the convergence test. The size of the marker is indicative of the total mass of a halo at the corresponding lookback time.

4.3 Results

this was done as a convergence test to ensure that our results do not change depending on the resolution. As can be seen (Figs. 4.2–4.9), DG15e9b-CT does not differ significantly from DG15e9b and any changes can be attributed to the stochastic nature of star formation, showing that the models are sufficiently converged. In DG10e9 and DG12e9, the resolution was increased because of the low star formation efficiency in low mass halos. The force resolution of each simulation is fixed by the mass of the gas particles and the star formation criteria: the softening length is the same as the smoothing length of a gas particle that satisfies the density threshold. These are given in column 7 in Table 4.1. The softening length is the same for all the baryonic and dark matter particles.

4.3 Results

In what follows, all simulated galaxy properties have been calculated at the present epoch, unless explicitly stated otherwise, to enable a comparison with dwarf galaxies observed in the nearby Universe. Some of the properties of the simulations at $z = 0$ are given in Table 4.2.

4.3.1 Baryonic Tully-Fisher relation

The baryonic Tully-Fisher relation (BTFR) relates the total baryonic mass to the circular velocity (see e.g. McGaugh, 2012). Since the inner mass density profile is strongly influenced by the gravitational coupling between gas and dark matter (Cloet-Osselaer *et al.*, 2012), a galaxy’s maximum circular velocity depends non-trivially on the total mass, including dark matter, and its star-formation history. Indeed, bursty star formation will be more efficient in reshaping the inner parts of the dark matter halo (Read & Gilmore, 2005, e.g.). The simulations including Pop III feedback yield dwarf galaxies that lie on the observed BTFR, in terms of total baryonic matter (Figure 4.2a) as well as the separate stellar and neutral gas component (Figure 4.2b and 4.2c, respectively). We confirm that these simulations reproduce the observed relation between stellar mass and neutral gas mass as well (Figure 4.2d). The control simulations without Pop III feedback fall outside the 3σ prediction interval of the regression lines fitted to the $v_c - M_{\text{bar}}$ and $v_c - M_{\star}$ observations: they produce ~ 10 times too many stars for their circular velocity. In short, our simulations including Pop III feedback succeed in creating faint, gas-dominated dwarf galaxies lying on the BTFR, such as Leo T, Leo P, and Pisces A, while the control simulations without Pop III feedback do not.

It is important to discuss the different measures for the rotational velocity of a galaxy here. McGaugh (2012) use both V_f , the velocity corresponding to the flat part of the rotation curve, and $W_{20}/2$. They find a much tighter BTFR using V_f

(1) Name	(2) M_* $10^6 \text{ M}_\odot$	(3) M_{HI} $10^6 \text{ M}_\odot$	(4) M_{halo} $10^9 \text{ M}_\odot$	(5) $M_{0.3}$ $10^7 \text{ M}_\odot$	(6) M_{half} $10^7 \text{ M}_\odot$	(7) M_V mag	(8) $B - V$ mag	(9) $V - I$ mag
DG9e9	0.10	1.14	4.94	2.65	0.77	-8.70	0.31	0.80
DG10e9	1.89	1.38	5.38	1.66	3.46	-10.98	0.54	1.00
DG12e9	0.15	0.26	6.33	1.98	0.71	-8.29	0.55	1.00
DG13e9a	0.25	3.68	9.53	1.39	0.84	-9.09	0.47	0.93
DG13e9b	1.23	3.07	5.93	0.58	2.06	-10.50	0.60	1.01
DG15e9a	3.31	1.73	9.75	3.52	2.16	-11.77	0.49	0.96
DG15e9b	4.41	26.35	9.22	1.20	2.67	-12.40	0.41	0.89
DG20e9	3.74	4.05	12.83	1.18	2.51	-11.80	0.52	0.99
DG50e9	14.83	66.40	34.70	0.58	5.97	-13.62	0.43	0.91
DG1e11	1187.60	462.08	63.61	3.27	53.94	-18.17	0.47	0.96
DG15e9b-CT	8.88	31.57	8.64	0.48	4.42	-12.98	0.45	0.93
DG10e9-NP3	20.43	41.47	7.04	0.24	10.54	-13.57	0.55	1.00
DG12e9-NP3	11.85	14.14	6.37	0.23	9.41	-12.94	0.62	1.01
DG13e9b-NP3	31.02	35.67	7.00	0.19	12.63	-14.00	0.60	1.01

(10) $\mu_{0,V}$ mag arcsec $^{-2}$	(11) R_e kpc	(12) $R_{30,V}$ kpc	(13) SFR $10^{-4} \text{ M}_\odot \text{ yr}^{-1}$	(14) v_c km s $^{-1}$	(15) σ_* km s $^{-1}$	(16) $\langle [\text{Fe}/\text{H}] \rangle_{\text{RGB}}$ dex	(17) $12 + \log(\text{O}/\text{H})$ dex	
DG9e9	24.52	0.14	0.43	1.07	14.54	10.9	-1.45	7.22
DG10e9	24.60	0.43	1.26	1.08	12.57	13.0	-1.46	7.94
DG12e9	25.46	0.18	0.45	0.14	11.92	12.0	-1.40	7.71
DG13e9a	25.29	0.24	0.63	0.43	19.72	7.8	-1.53	7.08
G13e9b	25.41	0.49	1.24	0.44	15.15	10.3	-1.66	7.74
DG15e9a	22.52	0.23	0.96	3.90	20.59	15.6	-1.23	8.16
DG15e9b	23.18	0.42	1.59	10.82	24.03	12.9	-1.44	7.64
DG20e9	23.68	0.40	1.40	4.32	14.73	12.3	-1.40	8.03
DG50e9	23.03	0.69	2.65	49.06	24.24	15.3	-1.48	7.93
DG1e11	19.32	1.02	5.99	1425.77	60.35	34.9	-0.91	8.76
DG15e9b-CT	23.64	0.69	2.39	11.12	19.92	11.6	-1.31	7.81
DG10e9-NP3	24.32	1.23	3.83	19.14	19.20	14.5	-1.60	8.11
DG12e9-NP3	25.01	1.26	3.46	5.62	15.16	14.2	-2.05	8.08
DG13e9b-NP3	23.89	1.23	4.12	20.22	19.59	15.3	-1.94	8.23

Table 4.2: Properties of the simulations at $z = 0$. (1): Name of the simulation, as described in Table 4.1, (2): The stellar mass, (3): the H I mass, (4): the virial mass of the dark matter halo, (5): total mass (baryonic and DM) within 300 pc, (6): total mass within the half light radius, (7): V-band magnitude, (8): $B - V$ colour, (9): $V - I$ colour, (10): central V-band surface brightness, (11): half light radius, (12): radius of the 30 mag arcsec $^{-2}$ isophote, (13): star formation rate, (14): circular velocity estimated from the H I linewidths ($v_c = W_{20}/2$), (15): stellar velocity dispersion, (16): mean stellar iron abundance, and (17): mean ionized gas oxygen abundance.

compared to using linewidths. The observations shown in Figure 4.2 use V_f while we used $W_{20}/2$ for the MORIA galaxies. This might explain the larger scatter for the simulations. We revisit this in section 6.2.

4.3.2 Cumulative star formation histories

In Figure 4.3, we present the CSFHs. The upper panel (Figure 4.3a) shows the CSFH of all the stars ever formed in the simulation. However, when we only

4.3 Results

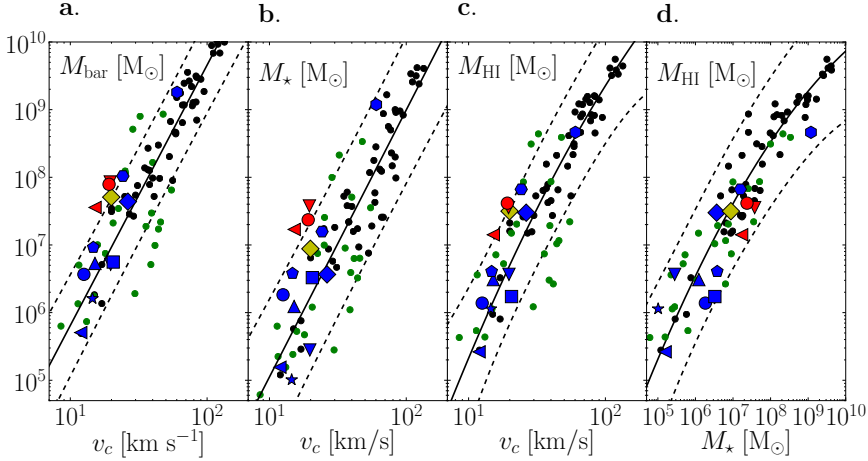


Figure 4.2: The baryonic Tully-Fisher relations (a, b, and c) and stellar mass versus H I mass (d). The MORIA simulations compared with the observations (McGaugh, 2012; Bernstein-Cooper *et al.*, 2014). v_c for the simulations are derived from their H I spectrum: $v_c = W_{20}/2$. The different symbols represent different simulations, as indicated in Table 4.1. Blue, red and yellow symbols are simulations with Pop III feedback, without Pop III feedback and the convergence test respectively, green symbols are symbols not discussed in detail, and black circles are observations. The solid lines are the regression lines through the data (linear in a and b, quadratic in c and d) and the dashed lines delimit the 3σ prediction interval.

consider the stars currently within a smaller projected distance from the galaxy center, the stellar mass assembly appears significantly more delayed. In Figure 4.3b, we present the CSFH of the stars currently within $1 R_e$, which is more in line with the area covered by current stellar-populations surveys (Weisz *et al.*, 2014). These more central CSFHs look strikingly like those of observed star-forming dwarf galaxies with comparable stellar masses in and near the Local Group (Weisz *et al.*, 2014). In contrast, previous simulations and semi-analytical models generally form too many stars early on (Weinmann *et al.*, 2012).

The cause for the radial dependence of the CSFH is that later stellar generations tend to be born within a more centrally concentrated volume than the earliest stars and that stars can migrate radially and become unbound and dispersed by the tidal forces accompanying merger events. Thus, older stars, which are subjected to more merger events than younger ones, are preferentially affected. In the turbulent ISM of the simulations with Pop III feedback, the first stars show

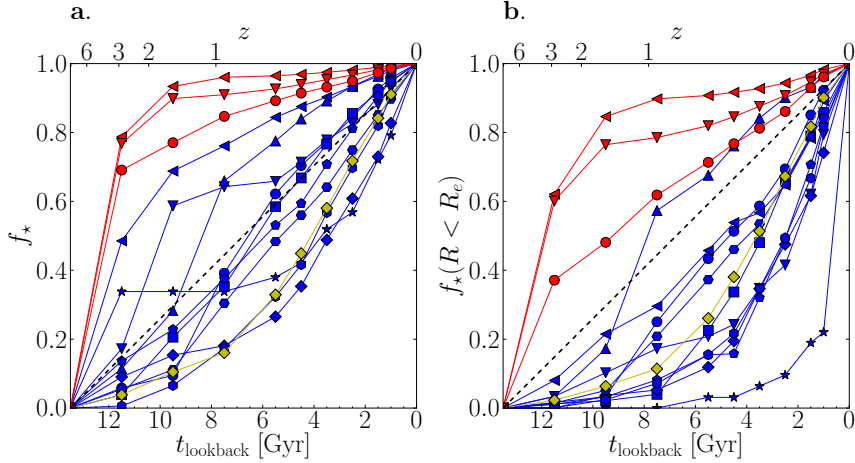


Figure 4.3: Cumulative star formation histories. The fraction of the stars formed by a certain lookback time is shown, derived from all stars (a) within the simulation, and (b) within $1 R_e$. Symbols and colours are the same as in Figure 4.2. The dotted lines shows the case for a constant star-formation rate.

significant outwards radial migration and are subsequently strongly affected by tidal forces. This is much less the case in the control simulations without Pop III feedback. As a consequence, these have much less radially varying CSFHs.

4.3.3 Metallicities

Metallicity measurements for nearby dwarf galaxies are often based on spectroscopic observations of bright stars on the Red Giant Branch (RGB) (e.g. Starkenburg *et al.*, 2010). The number of Pop II RGB stars within the half-light radius of the simulated dwarfs ranges from ~ 65 for the faintest one up to ~ 2000 for the brightest ones. The estimated number of Pop III RGB stars is zero in all simulated dwarfs presented here. The mean stellar iron abundance of stars that are now on the RGB in the simulated dwarfs agrees very well with the observations (Figure 4.4a). The oxygen abundance of the ionized gas surrounding star-forming regions is somewhat on the high side but also in broad agreement with the observations (Figure 4.4b). Oxygen is an element forged mostly by core-collapse supernovae while iron is formed abundantly in SNIa. The simulations without Pop III feedback form most of their stars early on and therefore have an inordinately large low-metallicity RGB population and hence fall significantly below the observed luminosity-metallicity relation of star-forming dwarf galaxies.

4.3 Results

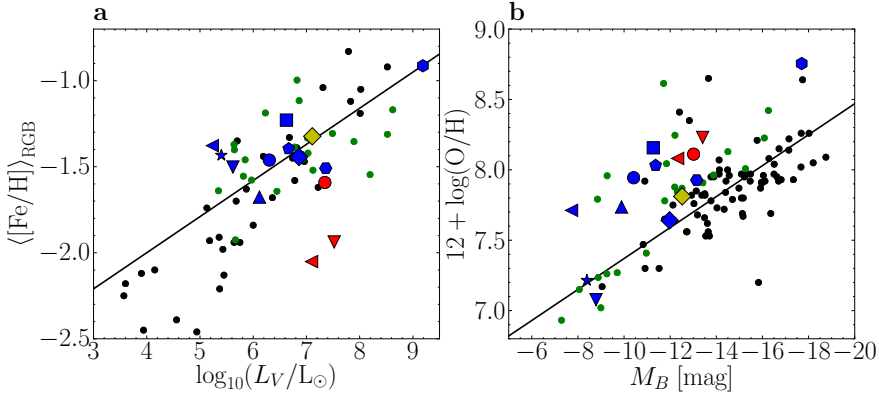


Figure 4.4: Metallicities. (a) The stellar Iron abundance and (b) the Oxygen abundance of the ionized gas around star forming regions of the simulations compared with observations of dwarf irregulars (Croxall *et al.*, 2009; Berg *et al.*, 2012; Skillman *et al.*, 2013; Kirby *et al.*, 2013). Symbols and colours are the same as in Figure 4.2. The fit to the data in (a) is $\langle [\text{Fe}/\text{H}] \rangle_{\text{RGB}} = -2.84 + 0.21 \log_{10}(L_V/L_\odot)$ (Kirby *et al.*, 2013) and in (b) $12 + \log(\text{O}/\text{H}) = 6.27 - 0.11 M_B / \text{mag}$ (Berg *et al.*, 2012).

4.3.4 Metallicity distribution functions

In the previous Section, we discussed the global metallicity of the simulated galaxies. However, metallicity distribution functions (MDFs) are available in the literature for many dwarf galaxies (e.g. Kirby *et al.*, 2013). To construct MDFs from our simulations, we again limit ourselves to the stars inside $1 R_e$ and take the predicted number of RGB stars within each stellar population into account. Figure 4.5 shows the MDF of Leo I, taken from Kirby *et al.* (2013), along with the MDFs of DG15e9b and DG12e9-NP3, which both have stellar masses similar to Leo I. The MDFs of DG12e9-NP3 and the other simulations without Pop III feedback all have a very long low-metallicity tail: $\sim 1\%$ of the RGB stars is in the form of stars with $[\text{Fe}/\text{H}] < -5$. This long tail is absent in observations and such metal-poor stars would be more easily detected. On the other hand, the low-metallicity tail is not present in DG15e9b and the other simulations including Pop III feedback and the MDF of DG15e9b shows very good agreement with that of Leo I. The simulation has slightly more metal-rich stars which could be because it is still actively star-forming, while Leo I has recently been quenched (Weisz *et al.*, 2014).

This gives further evidence that the first generation of stars did indeed have very

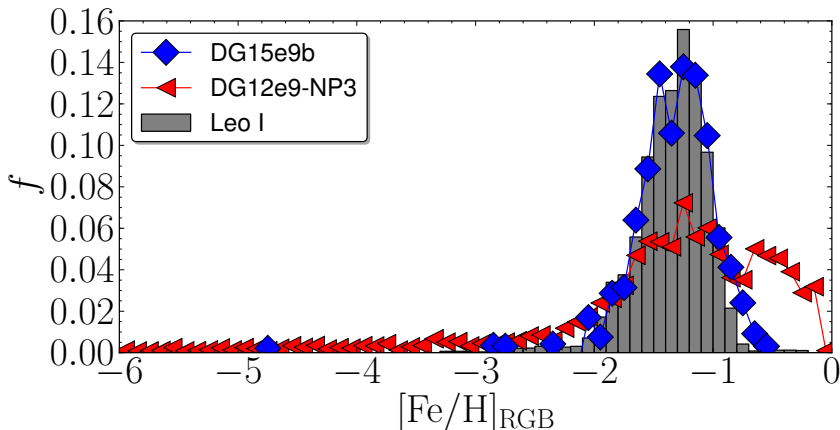


Figure 4.5: Metallicity distribution function. The MDFs of a simulation without Pop III feedback and of one including it, compared to the MDF of Leo I.

different properties than the stars we see today: if such extremely metal-poor stars would have an IMF similar to Pop II stars, they should be detectable. However to date, no such stars have been found in dwarfs (Starkenburg *et al.*, 2010).

4.3.5 $M_{\star} - M_{\text{halo}}$ -relation

The $M_{\star} - M_{\text{halo}}$ -relation relates the stellar mass to the dark matter mass and is typically obtained using the abundance matching technique; it is observationally less tractable but theoretically more straightforwardly calculated. We find that the models with Pop III feedback lie within the range predicted by abundance matching techniques (Guo *et al.*, 2010; Moster *et al.*, 2013; Behroozi *et al.*, 2013; Brook *et al.*, 2014, respectively the black, magenta, cyan, and gray line in Figure 4.6). Note that the simulations with the largest circular velocity do not necessarily have the most massive dark matter halos since the former also depends on the spatial distribution of the matter. At a given halo mass, the scatter on the stellar mass is considerable. This suggests that the abundance matching approach likely loses its applicability in this mass regime. The control simulations without Pop III feedback lie significantly above the predicted $M_{\star} - M_{\text{halo}}$ -relations. They show less scatter than the simulations with Pop III feedback because of the large amount of stars formed at high redshift. This will be less affected by the stochasticity of star formation. Star formation at low redshift is more stochastic but also happens at a much lower rate. Thus it will not alter the position in the

4.3 Results

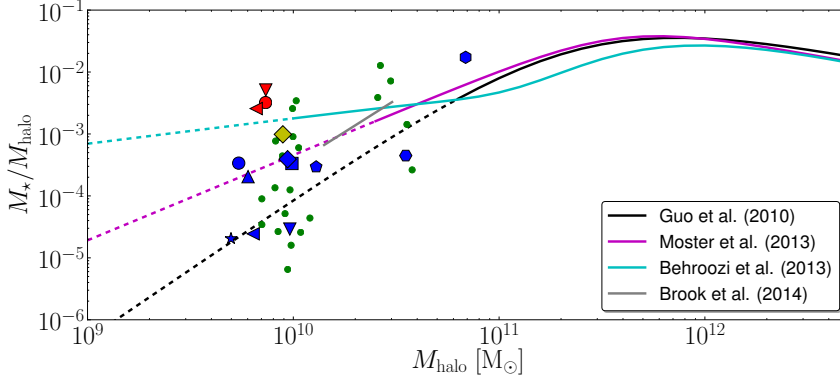


Figure 4.6: $M_{\star} - M_{\text{halo}}$ -relation. The simulations compared to the abundance matching relations of Guo *et al.* (2010) (black line), Moster *et al.* (2013) (magenta line), Behroozi *et al.* (2013) (cyan line), and Brook *et al.* (2014) (gray line). The observational scatter on the constrained parts of the relations (solid lines) is typically ~ 0.3 dex. The dashed segments of the lines indicate extrapolations to lower halo masses. Symbols and colours of the simulations are the same as in Figure 4.2.

$M_{\star} - M_{\text{halo}}$ relation much.

4.3.6 HI distribution

Figure 4.7 shows the H I distribution, rendered with the publicly available ray tracing package Splotch¹, overplotted with their H I density contours, with beam sizes of 40 pc. While this shows that our simulations look qualitatively like real dwarf galaxies, we can look at them more quantitatively.

The substructure in the ISM can be quantified using Fourier transform power spectra of the H I maps. These can be compared with radio observations of dwarf galaxies via the spectral index β of the power spectrum $P(k) \propto k^{\beta}$, where k is the wave number (Figure 4.8). Like for most real dwarfs in this luminosity regime, this index scatters around $\beta = -11/3 \approx -3.67$, the value expected in the case of a three-dimensional Kolmogorov-type incompressible, subsonic turbulence (Zhang *et al.*, 2012). This constitutes a test of the feedback prescription, responsible for blowing “holes” in the neutral gas and keeping the ISM thick and turbulent.

¹<http://www.mpa-garching.mpg.de/~kdolag/Splotch>

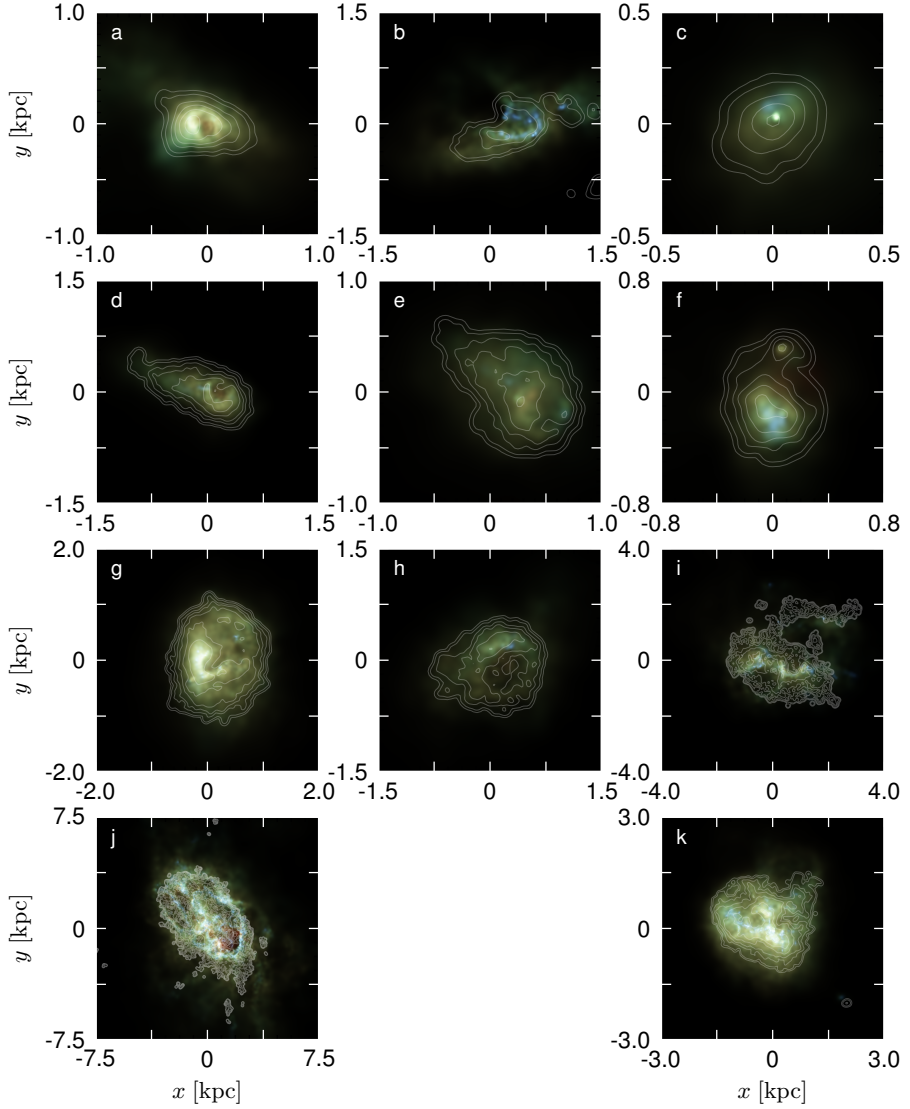


Figure 4.7: H I distribution of the simulations. The gas is rendered using the publicly available ray tracing software Splotch, overplotted with the H I density isophotes, with the lowest level $N_{\text{HI}} = 10^{19} \text{ cm}^{-2}$ with each next level an increment of a factor 4. The beam size for the H I contours is 40 pc. (a) DG9e9, (b) DG10e9, (c) DG12e9, (d) DG13e9L, (e) DG13e9E, (f) DG15e9L, (g) DG15e9E, (h) DG20e9, (i) DG50e9, (j) DG1e11 and (k) DG15e9b-CT.

4.3 Results

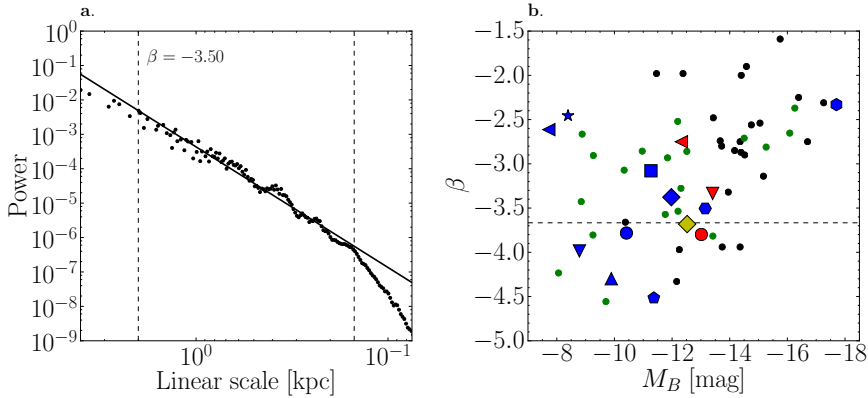


Figure 4.8: Spatial power spectra of the H I. (a) The power spectrum of the spatial distribution of H I in DG50e9. The solid grey line shows the power-law fit while the vertical dashed lines show the fitting range. (b) Spectral indices of the power spectra of our simulations, compared with dwarfs in the Little Things survey (Zhang *et al.*, 2012). The horizontal dashed line shows the value of 3D Kolmogorov-type turbulence. Symbols and colours are the same as in Figure 4.2.

4.3.7 Other properties

We also compare the optical colours, half-light radius, the central surface brightness, the central stellar velocity dispersion, the star-formation rate, the Kennicutt-Schmidt relation and the total mass within 300 pc and $1 R_e$ of our models with the observations (Figure 4.9a–i). We generally find good agreement between the observed scaling relations and our simulations with Pop III feedback. The control simulations without Pop III feedback are generally more extended and have lower central stellar and mass densities than observed dwarfs.

We note that we reproduce both $M_{0.3}$, the observed total mass within 300 pc (Strigari *et al.*, 2008, Figure 4.9h), and M_{half} , the total mass within the half-light radius (Collins *et al.*, 2014, Figure 4.9i). Furthermore, although these relations were determined for dSphs, with $L_V \lesssim 10^7 L_\odot$, our simulations agree with them over the entire luminosity range, and they predict an extension of this relation to $L_V \lesssim 10^9 L_\odot$. For some simulations, $M_{0.3}$ seems to be on the high side however. This is further discussed in the next Section.

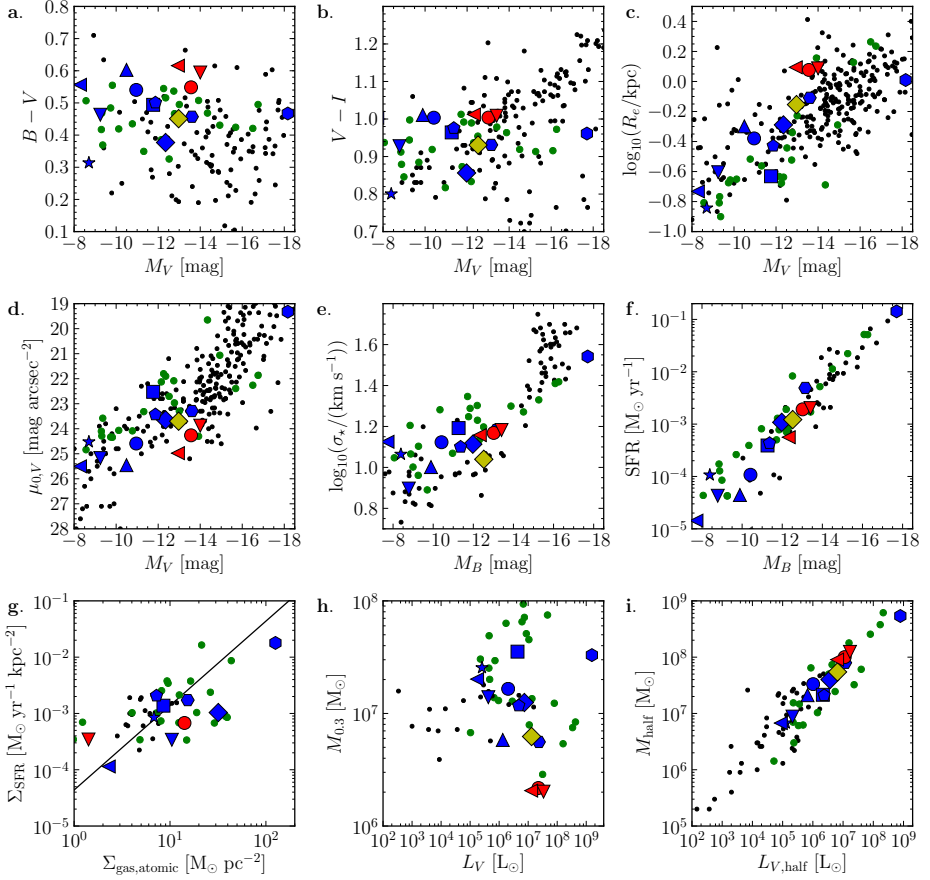


Figure 4.9: Other scaling relations. (a) $B - V$ colour, (b) $V - I$ colour, (c) Half light radius R_e , (d) Central surface brightness in the V-band $\mu_{0,V}$, (e) Stellar velocity dispersion σ_* , (f) Star formation rate, (g) Kennicutt-Schmidt relation, (h) Total mass within 300 pc, (i) Total mass within $1 R_e$ compared to observations (in grey; Graham *et al.*, 2003; Hunter & Elmegreen, 2006; Strigari *et al.*, 2008; Vennik & Hopp, 2008; de Rijcke *et al.*, 2009; Makarova *et al.*, 2009; McConnachie, 2012; Rhode *et al.*, 2013; McQuinn *et al.*, 2013; Karachentsev & Kaisina, 2013; Collins *et al.*, 2014; Hopp & Vennik, 2014; Tollerud *et al.*, 2015; Roychowdhury *et al.*, 2015). Symbols and colours are the same as in Figure 4.2.

4.3 Results

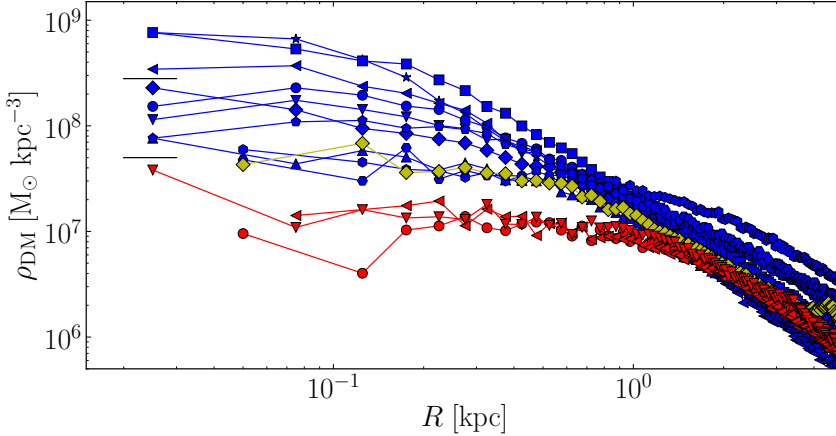


Figure 4.10: Dark matter density profiles of the simulations. Symbols and colours are as in Figure 4.2. Horizontal lines indicate the division between the different groups discussed in section 4.3.8.

4.3.8 Cusp versus core

Cosmological simulations including only gravity predict a universal density profile with a cusp in the central regions (Navarro *et al.*, 1996a). On the other hand, observations generally favor a cored density profile (de Blok *et al.*, 2008; Salucci *et al.*, 2012). It has been shown that including baryonic effects in simulations can drastically affect the dark matter density profiles and lead to cored density profiles (Read & Gilmore, 2005; Governato *et al.*, 2010; Cloet-Osselaer *et al.*, 2012).

Figure 4.10 shows the dark matter density profiles of the simulations. The models can be divided into three groups: DG9e9, DG12e9 and DG15e9a (group 1) have a high dark matter density in the central regions ($\rho_{\text{DM},0} \sim 10^9 \text{ M}_{\odot} \text{ kpc}^{-3}$); group 2 consists of the models with intermediary densities ($\rho_{\text{DM},0} \sim 10^8 \text{ M}_{\odot} \text{ kpc}^{-3}$); the central densities of the models without Pop III feedback (DG10e9-NP3, DG12e9-NP3 and DG13e9b-NP3, in red), who form group 3. We can now have a look to see if these groups manifest themselves in different scaling relations as well. From panels a and c in Figure 4.2, one might argue that group 1 lies lower on the BTFR, in terms of M_{bar} and M_{HI} , although only marginally. However, this is not true in terms of $M_{\star}$ (panel b). Group 3 lies significantly above the BTFR, indicating that they have unrealistic shallow inner density profiles.

Group 3 also has very early SFHs, although this will more likely be the cause of the shallow density profiles, rather than the other way around. Group 1 and

group 2 can not be distinguished from their SFHs.

Group 1 lies higher in the stellar metallicity plot (Figure 4.4a). One possible explanation is that because of their higher central densities, they can retain their metals more effectively. However, this is not necessarily the only explanation, since the division in groups based on their density profiles was done at $z = 0$ and it is not clear how much these density profiles change over time.

There is no clear distinction between group 1 and 2 in the $M_\star - M_{\text{halo}}$ -relation (Figure 4.6), indicating that the inner dark matter density profile and the total halo mass influence galaxy properties in different ways.

Not suprisingly, group 1 clearly has smaller R_e and $\mu_{0,V}$ (Figure 4.9c-d), while group 3 has larger R_e and $\mu_{0,V}$. Group 1 and group 3 also have larger, respectively smaller, $M_{0.3}$ (Figure 4.9h). DG1e11, which was categorized in group 2, also has a small R_e and large $M_{0.3}$, but this is mostly because of its high concentration of stars in its central region, rather than its dark matter properties. Group 1 seems to have a higher stellar velocity dispersion $\sigma_\star$, although only very slightly, since $\sigma_\star$ was determined at $1 R_e$, so at smaller radii for group 1. Because of the interplay between R_e and central density, all groups behave in the same way in terms of M_{half} (Figure 4.9i).

All of this indicates that to identify galaxies with a cusped dark matter density profile, the most likely candidates are the ones with small half-light radii and high total masses within a small, fixed radius. Furthermore, from our simulations it seems unlikely that any dSph galaxies analyzed in Strigari *et al.* (2008) have a cusped density profile.

4.3.9 Pop III to Pop II transition

The transition of Pop III to Pop II star formation occurs, in terms of metallicity, at $[\text{Fe}/\text{H}] = -5$ in our models. However, it is not immediately clear at what time this transition occurs and how abrupt it is. Figure 4.11 shows the star formation rate of Pop III stars and of Pop II stars (in red and blue, respectively) for one of our least massive (DG10e9, top panel) and our most massive model (DG1e11, bottom panel). The reason we show DG10e9 rather than DG9e9, is because there is no star formation in this model for $3 \lesssim z \lesssim 1$, as seen in Figure 4.3. Before 1 Gyr, the Pop III stars dominated the overall star formation, although there is already some low level Pop II star formation this early on. This Pop III star formation declines in time, after which the Pop II star formation starts dominating the global star formation and after some time, all stars are being formed out of sufficiently enriched gas. The exact time of the last Pop III stars being formed seems to differ significantly between the two models shown in Figure 4.11: the transition occurs much earlier for the least massive model, at $t \approx 1.6$ Gyr, while the most massive model still has Pop III star formation until $t \approx 3.5$ Gyr.

4.4 Summary

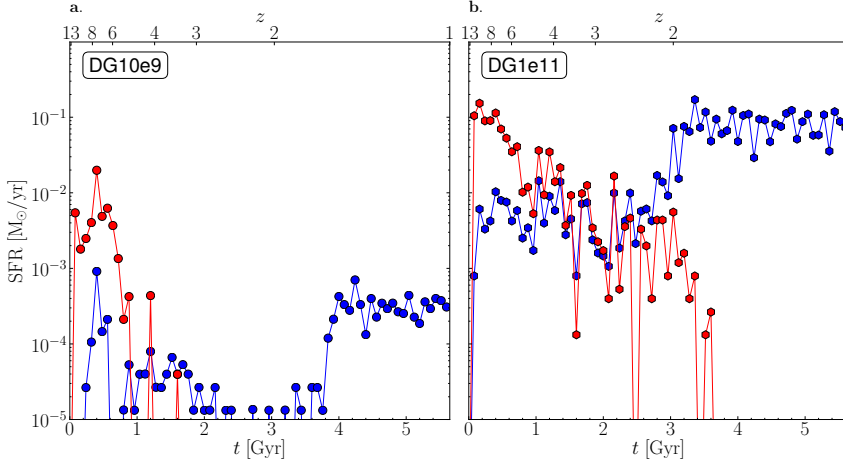


Figure 4.11: Transition of Pop III to Pop II star formation. Star formation rate of the Pop III stellar populations (red) and of the Pop II populations (blue) of DG10e9 (a) and DG1e11 (b), one of our least and our most massive models, respectively. All stellar particles formed in the simulations are taken into account.

This trend for later transition times for more massive galaxies is true for all our simulations.

Although this gives a general idea of how and when this transition between star formation modes might have occurred in the universe, we do not wish to over-interpret these results, since the cut-off metallicity is not well constrained and the transition will most likely have been more smoothly than the sharp bimodal model we obtained in our models. However, it does serve as an important check and we confirm that we do not have any Pop III star formation at low redshift, which would be more easily detected observationally.

4.4 Summary

We have shown that, with the inclusion of the fierce UV radiation emitted by Pop III stars in numerical simulations, we can reproduce the most important observed properties of isolated gas-rich dwarf galaxies within the Λ CDM paradigm. This energetic feedback from the very first generation of stars is crucial in suppressing the star-formation rate in a dwarf's progenitors and thus delaying its star-formation history. We stress that no parameters were fine-tuned to obtain

these results.

Furthermore, the expected number of Pop III RGB stars is essentially zero, in agreement with the most metal-poor stars found in dwarf galaxies (Starkenburg *et al.*, 2010).

Of course, there is still room for improvement. For instance, the Oxygen abundances predicted by our simulations are generally on the high side, as is the scatter between the values for the total mass within 300 pc. It remains to be seen if and how these remaining disagreements between models and data can be resolved.

In short, we have provided numerical evidence that the first stars that formed in the early Universe are essential for producing isolated, gas-dominated dwarf galaxies in simulations with broadly the same chemical, kinematical, and structural properties as real dwarf systems in the nearby Universe.

Finally, we wish to stress the crucial importance of mimicking the observations as closely as possible (e.g. by deriving RGB-weighted mean metallicities). This is the only meaningful way in which the validity of numerical simulations can be tested.

Too Big To Fail is in the Eye of the Beholder

There's no I in collaboraton.

Ken M

R. Verbeke , E. Papastergis, A.A. Ponomareva, S. Rath and S. De
Rijcke (2017)

A&A submitted

*"A new astrophysical solution to the Too Big To Fail problem -
Insights from the MoRIA simulations"*

5.1 Introduction

IN this Chapter, we investigate the TBTF problem for field dwarfs. We take to heart the message from P16: if Λ CDM is correct, then late-type field dwarfs should have higher circular velocities than is estimated from their H I kinematics. In order to investigate such a possible mismatch between the maximum circular velocity as inferred from gas kinematics and its actual value, we perform H I observations of a set of simulated dwarf galaxies. In Section 5.2, we briefly present the MoRIA simulations and the procedure to construct and analyse mock H I data-cubes. In Section 5.3, we fit a halo profile to the outermost datapoint of the rotation curves of the simulated galaxies and compare with the results of P16. In Section 5.4 we give some possible explanation for these results. Our conclusions are presented in Section 5.5.

For clarity, we define the different types of velocities used throughout this Chapter here:

- $v_{\text{rot,HI}}(R)$
the mean tangential velocity of the H I gas at a radius R from the galaxy

center. This can be determined from observations by fitting a tilted-ring model to the H I velocity field or the full data-cube.

- $v_{\text{circ}}^{\text{obs}}(R)$
the circular velocity derived from the $v_{\text{rot,HI}}(R)$ profile by correcting for asymmetric drift (see Section 5.2.4).
- $v_h^{\text{true}}(R)$
the “true” circular velocity profile, inferred from the total enclosed mass profile $M(R)$ as

$$v_h^{\text{true}}(R) = \sqrt{\frac{GM(R)}{R}}. \quad (5.1)$$

- $v_{\text{out,HI}} = v_{\text{circ}}^{\text{obs}}(R_{\text{out}})$
the outermost value of the rotation curve.
- $v_{h,\text{max}}^{\text{true}} = \max(v_h^{\text{true}})$
the maximum circular halo velocity.
- $v_{h,\text{max}}^{\text{fit}}, v_{h,\text{max}}^{\text{NFW}}, \text{ or } v_{h,\text{max}}^{\text{DC14}}$
the maximum circular velocity obtained by fitting an NFW or DC14 profile to $v_{\text{out,HI}}$. Denoted by $v_{h,\text{max}}^{\text{fit}}$ in general and $v_{h,\text{max}}^{\text{NFW}}$ or $v_{h,\text{max}}^{\text{DC14}}$ when the halo profile is specified.
- W_{50}
The FWHM of the galactic 21cm emission line profile, corrected for inclination to an edge-on view.

All except for W_{50} refer to a spatially resolved kinematic measurement or calculation. W_{50} on the other hand is derived from the spatially unresolved H I spectrum. Since W_{50} does not correspond to any specific radius, it does not generally contain enough information to estimate the mass of the host halo by fitting a certain mass profile. However, W_{50} measurements exists for large samples of galaxies, which allows for an accurate measurement of the number density of galaxies as as function of W_{50} , i.e. the VF.

5.2 The MoRIA simulations

We again use the MoRIA suite of N-body/SPH simulations of late-type isolated dwarf galaxies. We again select 10 simulated galaxies to discuss in more detail in this chapter. An overview of some of the basic properties of the 10 selected MoRIA dwarfs is presented in Table 5.1. M-1 to M-5 (M-6 to M-10) have a mass resolution of $4230 M_{\odot}$ ($10515 M_{\odot}$) for its baryonic component and a force

5.2 The MORIA simulations

(1)	(2)	(3)	(4)	(5)	(6)	(7)	(8)	(9)	(10)	(11)	(12)
Name	Symbol	$\log_{10}(M_*)$ [$M_\odot$]	$\log_{10}(M_{\text{HI}})$ [$M_\odot$]	$\log_{10}(M_{200})$ [$M_\odot$]	M_V [mag]	q	R_{out} [kpc]	$v_{\text{out, HI}}$ [km s $^{-1}$]	$v_{h, \text{max}}^{\text{true}}$ [km s $^{-1}$]	$W_{50}/2$ [km s $^{-1}$]	σ_* [km s $^{-1}$]
M-1	◇	6.56	7.48	9.97	-12.35	0.49	1.68	21.66	31.04	8.34	10.76
M-2	★	6.59	7.29	9.95	-12.52	0.63	1.21	39.87	32.83	6.15	17.86
M-3	○	6.87	7.31	9.93	-12.77	0.74	1.13	19.45	28.04	6.31	12.60
M-4	△	7.41	7.83	10.00	-13.92	0.41	2.37	17.04	41.42	11.72	12.46
M-5	▽	7.55	7.34	10.01	-14.35	0.53	1.09	41.69	30.36	10.03	18.77
M-6	◁	7.71	7.94	10.55	-14.93	0.76	1.93	39.04	36.61	12.73	21.34
M-7	▷	8.00	8.49	10.41	-15.69	0.58	4.18	36.31	36.16	14.30	19.68
M-8	□	8.33	8.64	10.47	-16.50	0.61	4.87	45.79	43.45	8.02	26.25
M-9	◊	8.53	8.59	10.42	-16.75	0.60	3.43	40.54	66.02	15.09	25.09
M-10	○	9.07	8.66	10.84	-18.17	0.56	4.15	53.94	67.30	38.27	33.84

Table 5.1: Properties of the 10 selected MORIA simulations at $z = 0$. (1) The name of the simulation, (2) the symbol used throughout the plots, (3) the stellar mass, (4) the H I mass, (5) the halo virial mass, (6) the total V-band magnitude, (7) the intrinsic flattening of the H I, (8) the H I radius, (9) the outermost value of the rotation curve, (10) the maximum circular velocity of the halo, (11) the half-width-half-max of the H I, and (12) the velocity dispersion of the stars at R_{out} .

resolution of 9.8 pc (13 pc). We note that these are different then the ones discussed in Chapter 4, although there is some overlap: DG15e9b and DG1e11 are the same as M-1 and M-10, respectively.

5.2.1 H I disk sizes and flattening

We aim to investigate H I rotation curves, with strong focus on the outer-most data point. It is therefore very important that the simulated dwarf galaxies have realistic H I disk sizes and shapes. In the previous chapter, we already showed that the MORIA dwarfs have an atomic ISM with realistic spatial substructure, as quantified by the H I power spectrum.

Here, we also investigate the flattening and size of the H I disks. For this, we produce H I surface density contour maps of the H I and fit ellipses to the contour corresponding to a column density of $\Sigma_{\text{HI}} = 1 M_\odot \text{ pc}^{-2} \approx 1.25 \times 10^{20} m_{\text{H}} \text{ cm}^{-2}$. We do this for different orientations and take the minimum value of the flattening q , defined as the ratio of the minor and major axis of the ellipse. In other words, q is the intrinsic axis ratio of the galaxy. The frequency distribution of the axis ratio q of the simulated MORIA galaxies is shown in Figure 5.1, along with that of observed dwarf galaxies, derived by Roychowdhury *et al.* (2010) for the FIGSS sample of faint galaxies. In Figure 5.2, we show the total H I mass, denoted by M_{HI} , in function of the disk size, R_{out} , both for simulated and observed galaxies. The disk size is defined as the major axis of the elliptical contour corresponding to a column density of $\Sigma_{\text{HI}} = 1 M_\odot \text{ pc}^{-2}$.

We generally find good agreement with the observed flattening distribution al-

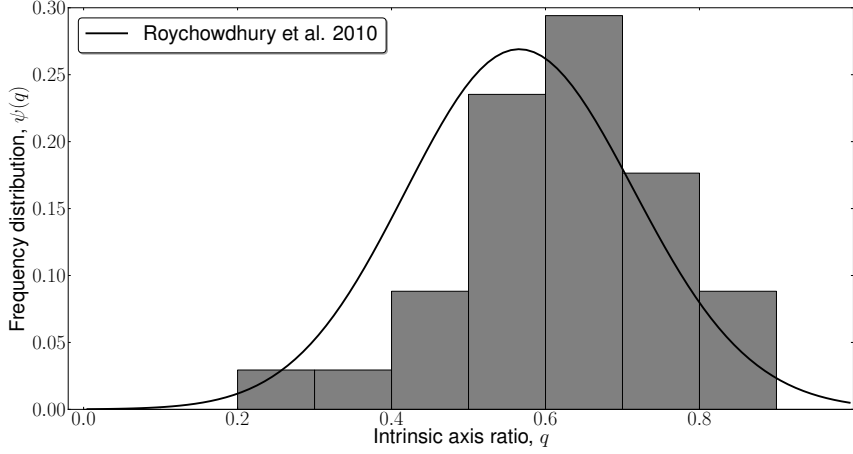


Figure 5.1: Histogram of the axis ratios of the MORIA dwarf galaxies versus the frequency distribution obtained by Roychowdhury *et al.* (2010) for the FIGGS galaxies.

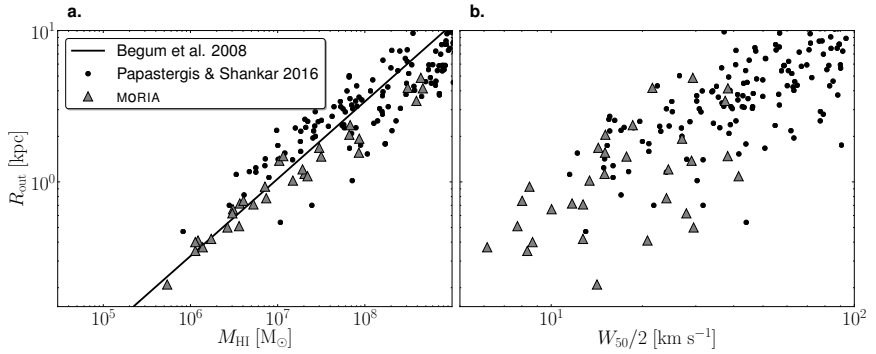


Figure 5.2: **a.** Atomic gas mass, M_{HI} versus H I disk size, R_{out} , and **b.** $W_{50}/2$ versus R_{out} of the MORIA dwarfs (in red) versus observations compiled in P16. For the simulations, R_{out} is the semi-major axis of the best fitted ellipse to the contour with $\Sigma_{\text{HI}} = 1 M_{\odot} \text{ pc}^{-2}$. The relation in panel **a.** is the one found for the FIGGS sample at $1 M_{\odot} \text{ pc}^{-2}$ ($\log(M_{\text{HI}}) = 1.96 \log(2R_{\text{out}}) + 6.37$; Begum *et al.*, 2008a).

5.2 The MoRIA simulations

though the MoRIA dwarfs appear to have slightly thicker H I disks than the observed dwarfs. However, the MoRIA dwarfs were not intended to be equivalent to the FIGGS sample. Indeed, most of the FIGGS galaxies have $M_{\text{HI}} \sim 10^7 - 10^9 M_{\odot}$ (Figure 1c in Begum *et al.*, 2008b) whereas more than half of the MoRIA dwarfs lie in the $M_{\text{HI}} \sim 10^6 - 10^7$ regime (see Figure 5.2). Roychowdhury *et al.* (2010) also note that galaxies with high inclinations may be overrepresented in their sample which might lead to a slight underestimate for the mean intrinsic axis ratio $\langle q \rangle$. Furthermore, they assumed in their analysis that the gas disks are oblate spheroids, and showed that $\langle q \rangle$ would be higher when assuming a prolate spheroid. Galaxies are not necessarily oblate spheroids (e.g. Cloet-Osselaer *et al.*, 2014), so the real $\langle q \rangle$ might be higher. Considering all this, it is remarkable that we find a distribution that looks so similar to the observed one.

As can be seen in Figure 5.2, the sizes of the H I disks of the MoRIA dwarfs are realistic as well: they follow the same mass-size relation and FWHM-size relation as the observed galaxies compiled in P16. This is of crucial importance because it determines the position of the outermost datapoint to which the circular velocity profile is fitted in order to estimate $v_{h,\text{max}}^{\text{fit}}$.

5.2.2 Mock data cubes

The procedure to produce a cube of 21cm data for a MoRIA dwarf is as follows. First, we tilt the galaxy such that its angular momentum vector is inclined by 45° with the line of sight. Then, the mass of each gas particle is assigned to a cell in a three dimensional grid based on its projected position and its line-of-sight velocity. The velocity grid is chosen with a resolution of 2.5 km s^{-1} . To account for thermal broadening, the H I mass of each gas particle is smeared out over neighboring velocity channels using a Gaussian with a dispersion given by

$$\sigma_{\text{TB}} = \sqrt{\frac{kT}{m_p}}, \quad (5.2)$$

where T is the temperature of the particle, k is the Boltzmann constant and m_p the proton mass. The gas is allowed to cool down to $T = 10 \text{ K}$ while it becomes fully ionized around $T \sim 10^4 \text{ K}$. So the thermal broadening achieves values in the interval $0.29 \text{ km s}^{-1} \lesssim \sigma_{\text{TB}} \lesssim 10 \text{ km s}^{-1}$. Finally, each velocity channel is convolved with a Gaussian beam profile as well. The full widths at half the maximum (FWHM) of the beams are shown in the top left panels in Figure 5.3. The beam size was chosen so that it fits at least 10 times within the HI radius of the galaxy. For the simulations presented here, this comes down to 100 pc for the ones with $R_{\text{out}} \sim 1 \text{ kpc}$ (M-1, M-2, M-3, M-5, and M-6) and 200 pc for the larger ones (M-4, M-7, M-8, M-9, and M-10). The resolution of the spatial grid

is chosen so that the beam size corresponds to 5 pixels. The three-dimensional mass grid is then saved in the FITS format.

5.2.3 Rotation curves

To achieve a realistic comparison analysis, we opt for two observational analysis codes to derive rotation curves for the MoRiA dwarfs based on their radio data cubes: GIPSY (The Groningen Image Processing SYstem; van der Hulst *et al.*, 1992) and ^{3D}Barolo (Di Teodoro & Fraternali, 2015). GIPSY has a built-in routine, ROTCUR, which fits a tilted-ring model to the H I velocity field (Begeman, 1989). Of the full suite of MoRiA dwarfs, we selected ten with velocity fields and shapes that are sufficiently relaxed to be amenable to analysis with ROTCUR. The ones that were not selected had a very irregular H I morphology or velocity field due to their low masses. ^{3D}Barolo fits a model directly to the full data-cube, which makes it useful for a comparison with the GIPSY results. The tilted-ring model in ^{3D}Barolo is populated with gas clouds at random spatial positions. This feature makes this code very useful for determining the kinematics of dwarf galaxies with sometimes highly disturbed gas distributions.

Ideally, given the way we produce the datacubes, one would expect the center of each ring in the tilted-ring model to coincide with the nominal galaxy center grid, its inclination to be 45°, and its position angle (PA) to be 90°. However, strongly disturbed and warped disks can lead to tilted-ring models with the apparent ring centers, inclinations, and PAs significantly shifted away from their expected values. The parameters are chosen by fitting an ellipse to the isodensity contour of $\Sigma_{\text{HI}} = 1 \text{ M}_{\odot} \text{ pc}^{-2}$. These were checked and adjusted so that, for instance, the rotation would be around the minor axis. The inclination is typically chosen to its true value of 45° unless the shape of the galaxy clearly implies a different inclination. The chosen ellipses are shown in the top panels of Figure 5.3. The adjusting of the parameters will typically lead to a smaller maximal radius than R_{out} . Also, the isodensity contours are not perfect ellipses (the chosen rings will thus go through areas with higher densities), leading to $\Sigma_{\text{HI}}(R_{\text{out}}) > 1 \text{ M}_{\odot} \text{ pc}^{-2}$. The systematic velocity is chosen as roughly the value of the center (typically close to 0 km s⁻¹). We keep these values fixed for each radius. The rotational velocity is thus the only parameter that is fitted.

5.2.4 Pressure support corrections

In the low-mass systems under investigation here, pressure support is expected to be significant, entailing a sizable correction. For the latter, we follow the approach typically used in observational studies of dwarf kinematics (e.g. Lelli

5.2 The MoRIA simulations

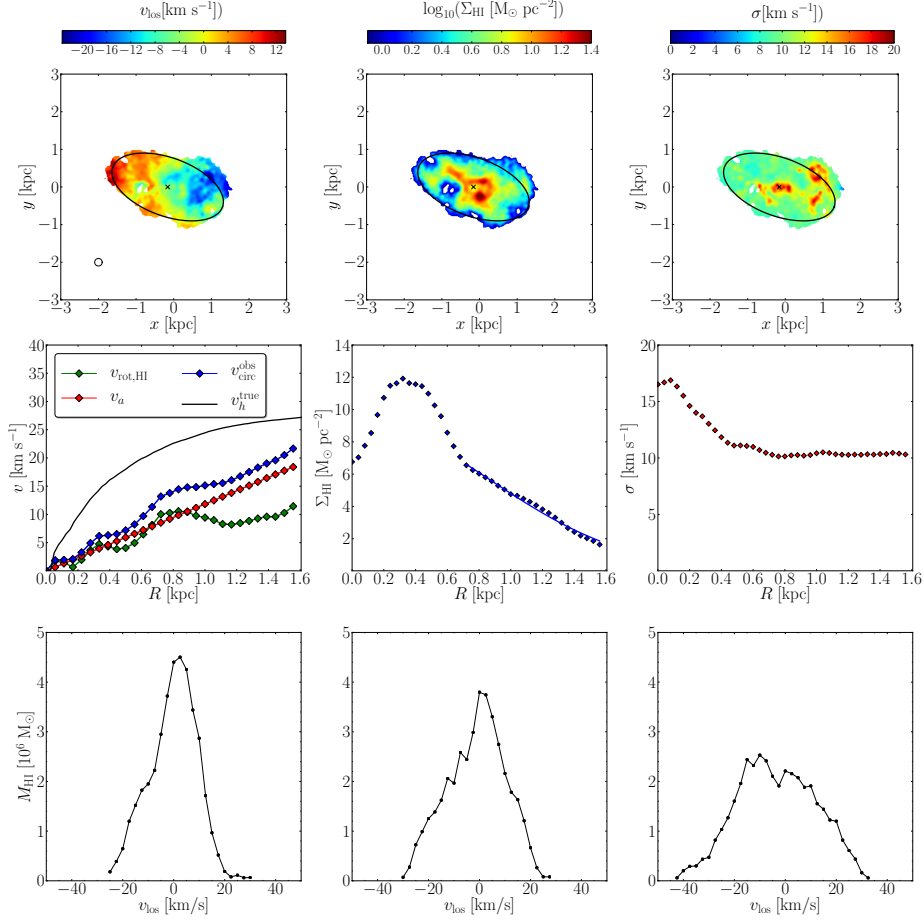


Figure 5.3: **Top panels:** velocity field (*left*), density (middle) and dispersion (right) map of M-1. The beam size used is shown in the top left panel. **Middle panels:** The rotation curves, with the rotational velocity obtained using tilted ring fitting in GIPSY in green, correction for pressure support in red, full circular velocity in blue and theoretical halo velocity in black (left), H I density profile, with the fit necessary for the pressure support correction plotted as a solid line shown over the area that was used for the fit (middle), H I velocity dispersion profile (right). Bottom panels: the global H I profile for the inclined view (left) as shown in the top panels and for two edge-on views (middle and right). The velocity bin width is the same as the channel width: 2.5 km s^{-1} .

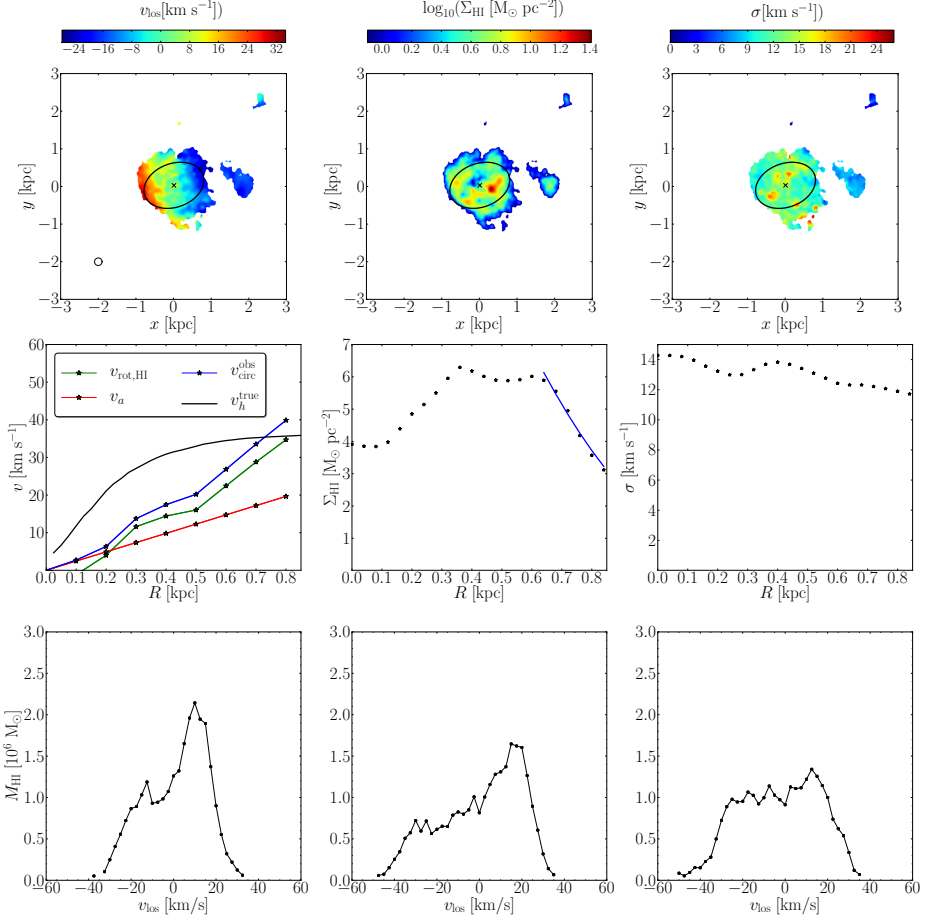


Figure 5.3: *Continued.* Overview of M-2.

5.2 The MoRIA simulations

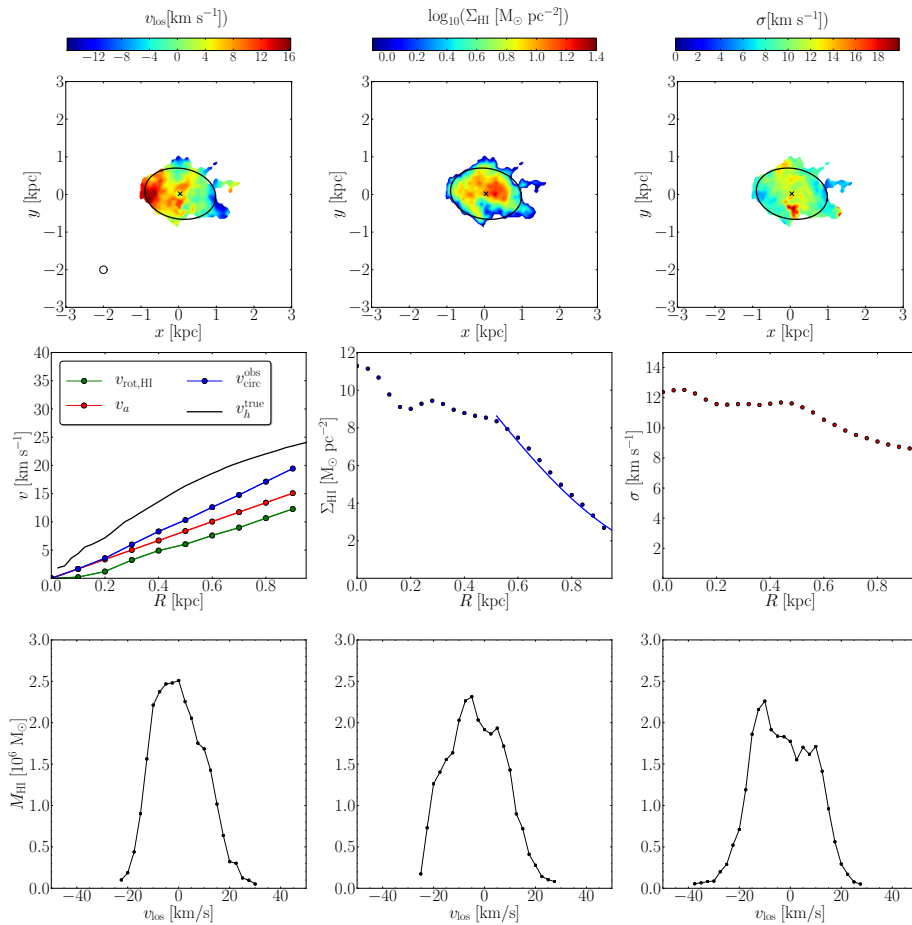


Figure 5.3: *Continued.* Overview of M-3.

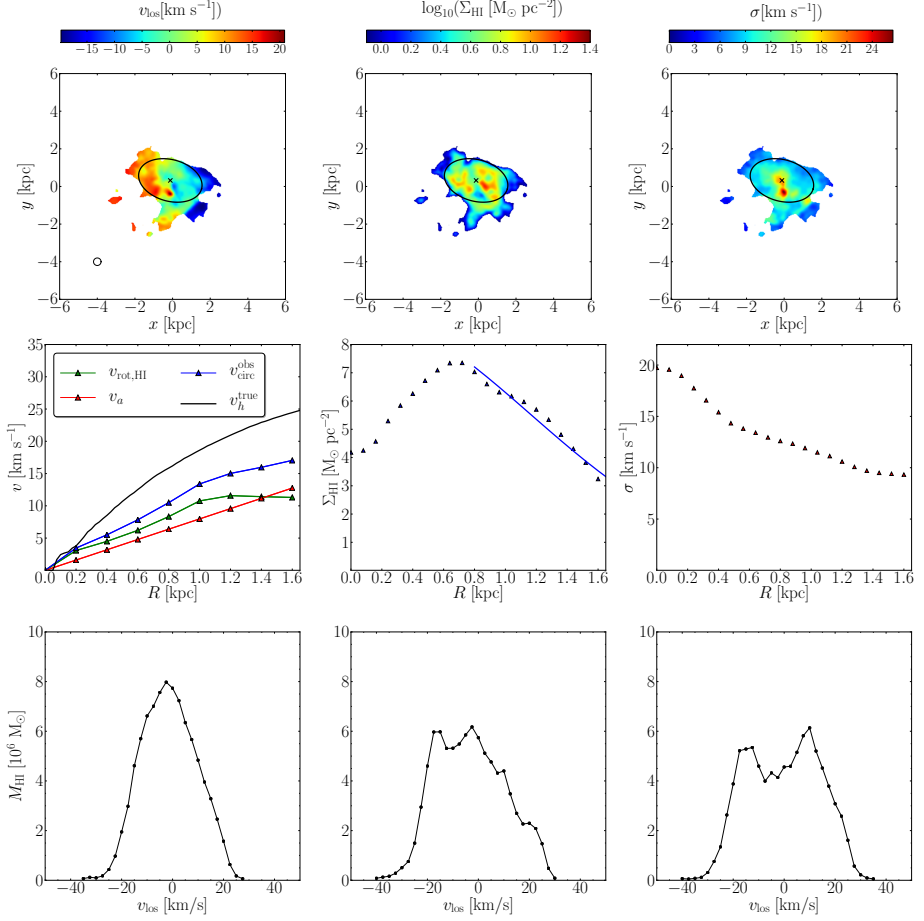


Figure 5.3: *Continued.* Overview of M-4.

5.2 The MoRIA simulations

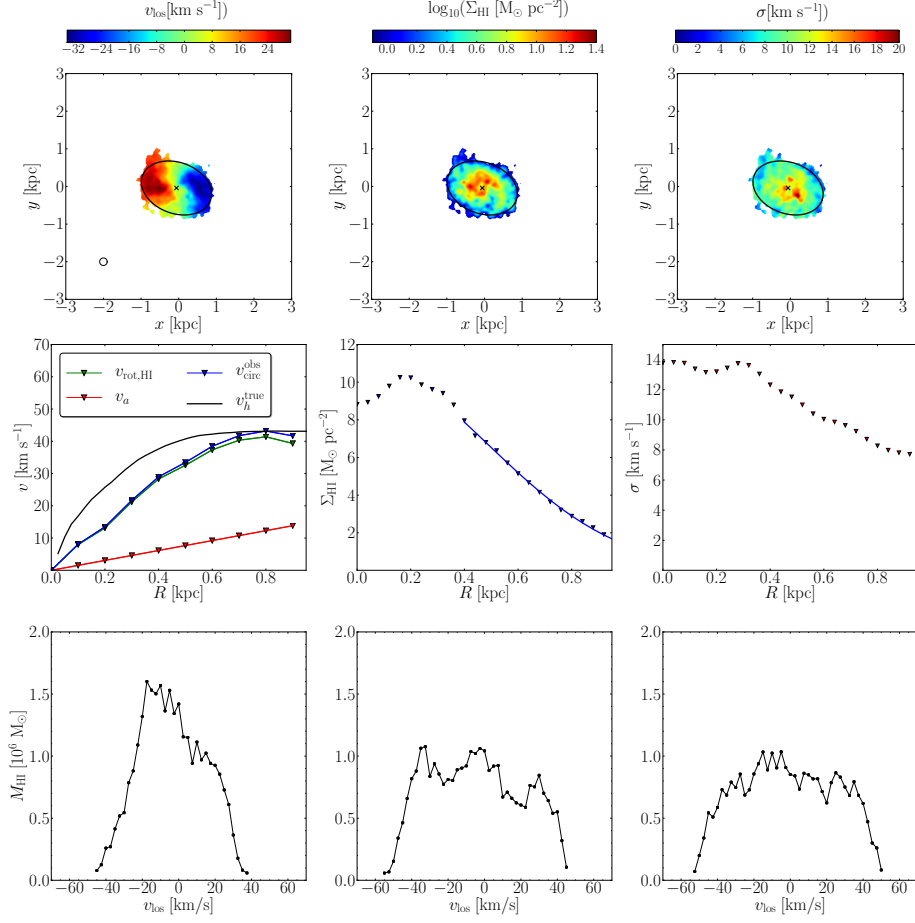


Figure 5.3: *Continued.* Overview of M-5.

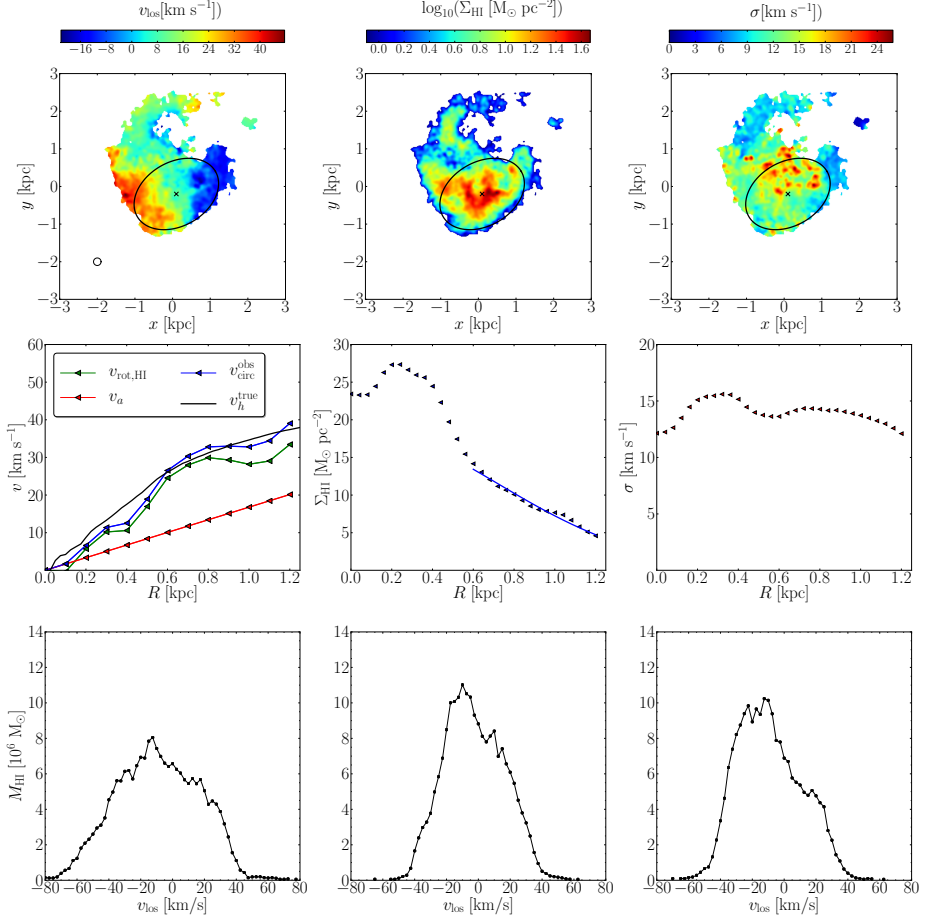


Figure 5.3: *Continued.* Overview of M-6.

5.2 The MoRIA simulations

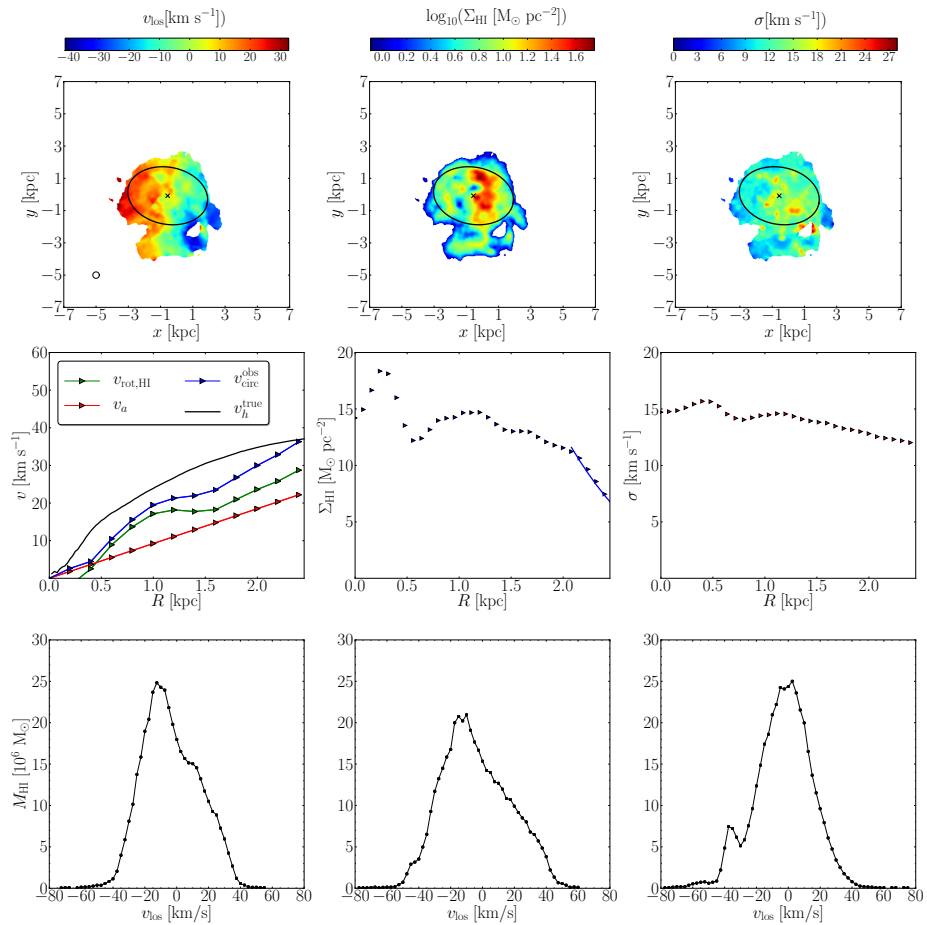


Figure 5.3: *Continued.* Overview of M-7.

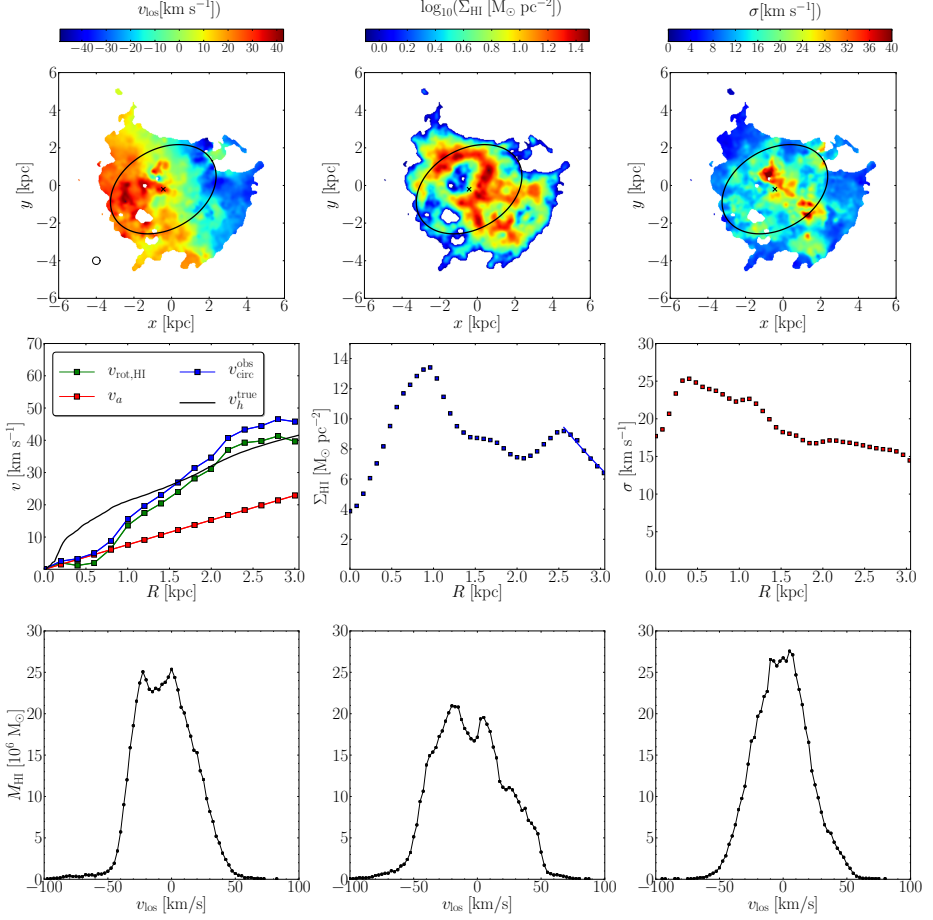


Figure 5.3: *Continued.* Overview of M-8.

5.2 The MoRIA simulations

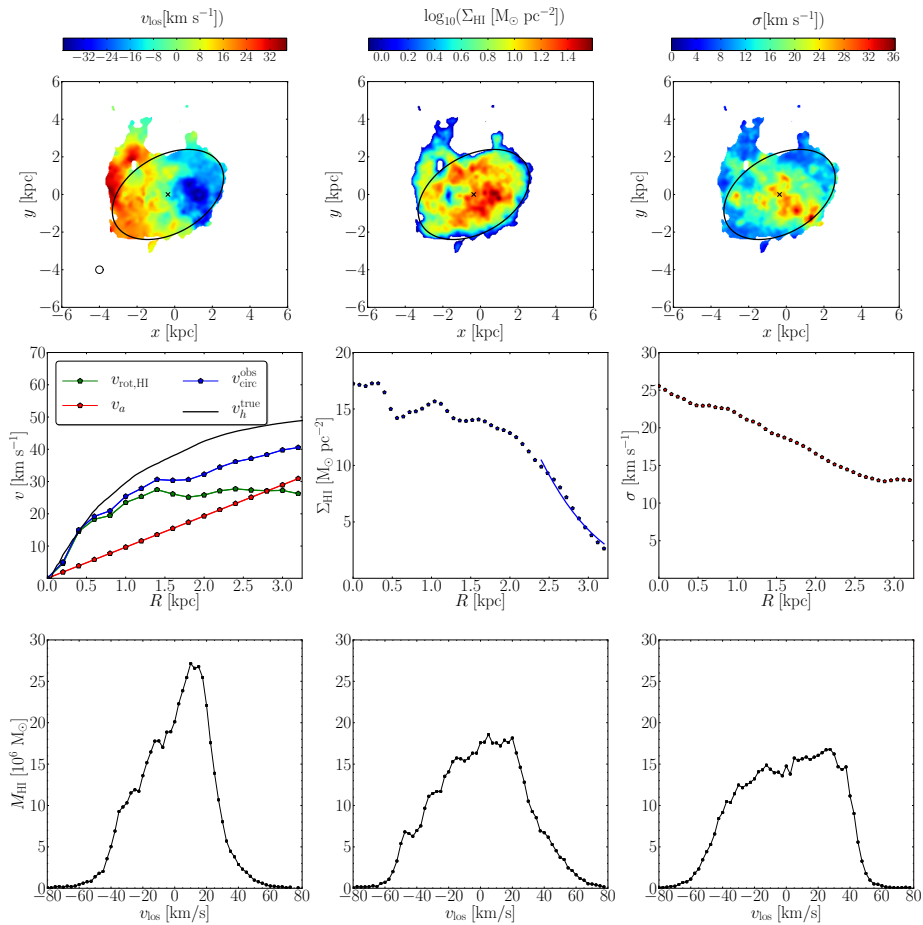


Figure 5.3: *Continued.* Overview of M-9.

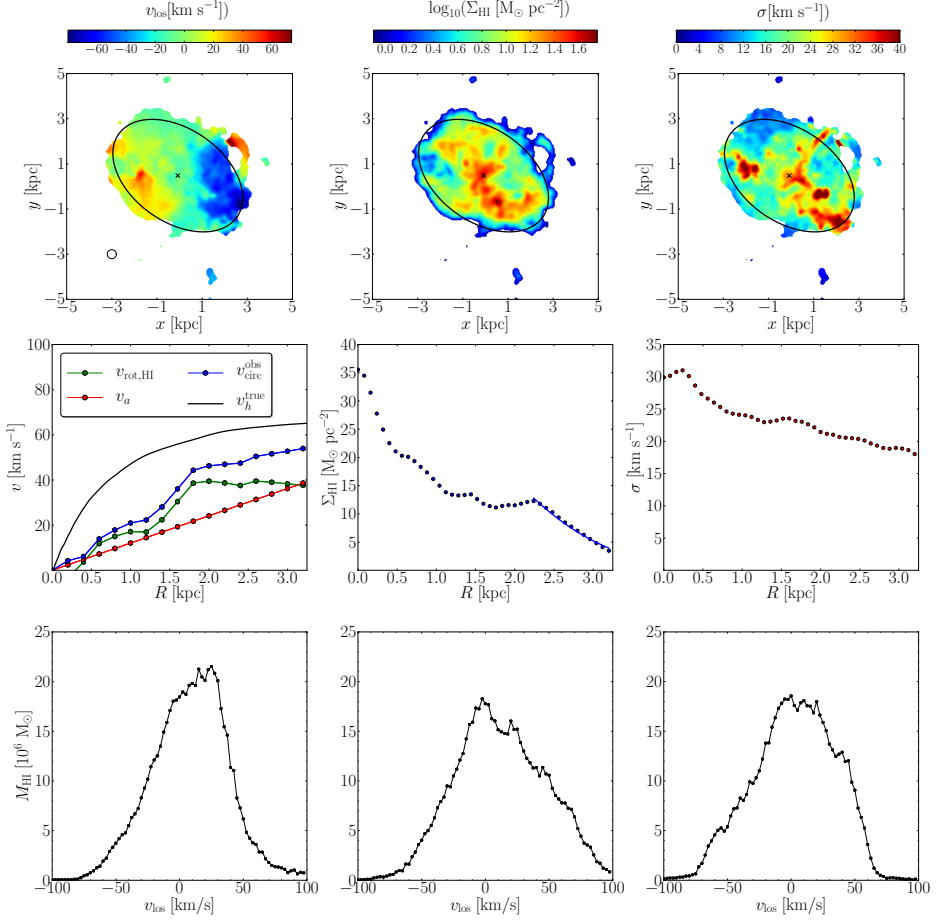


Figure 5.3: *Continued.* Overview of M-10.

5.2 The MoRIA simulations

et al., 2012b). The pressure support correction is given by

$$v_a^2(R) = -\sigma^2 \frac{\partial \ln(\sigma^2 \Sigma_{\text{HI}})}{\partial \ln R}, \quad (5.3)$$

where σ is the velocity dispersion and Σ_{HI} is the intrinsic gas surface density. We assume a prescription for Σ_{HI} of the form

$$\Sigma_{\text{HI}}(R) = \Sigma_0 \exp(-R^2/2s^2), \quad (5.4)$$

with s a radial scale length. Since we are only interested in v_{out} , the fit is performed for the outer regions and the velocity dispersion is assumed constant at the value at the outer edge of the galaxy. This is justified since the radial variation of σ is typically small. The pressure support correction then becomes

$$v_a^2(R) = \sigma_{\text{out}}^2 \frac{R^2}{s^2}. \quad (5.5)$$

We have also tried other commonly used prescriptions for Σ_{HI} , as well as for $\sigma^2 \Sigma_{\text{HI}}$ (e.g. Oh *et al.*, 2015). This led to the same general results. Given the rotation velocity v_{rot} provided by the tilted-ring fit and the pressure support correction v_a , we obtain an estimate for the true circular velocity as

$$v_{\text{circ}}^{\text{obs}}(R) = \sqrt{v_{\text{rot,HI}}^2(R) + v_a^2(R)}. \quad (5.6)$$

The obtained rotation curves are shown in the lower left panels of Figure 5.3. In this Figure, we show the three first moment maps of the data cube, the rotation and circular velocity curves, the radial H I density profile, and the H I velocity dispersion profile. In the density profile diagram, the region where the parameters of the pressure support model have been determined is indicated. In this region, the pressure support correction can be computed relatively reliably; at smaller radii, the resulting pressure support correction should be taken with a grain of salt.

In Figure 5.4, we compare rotation curves determined with GIPSY and with ^{3D}Barolo. For the tilted-ring analysis with ^{3D}Barolo, we used 20 rings, each with a radial size of 10arcsec $\approx$ 50parsec. We keep all parameters the same as in the analysis with GIPSY, with the exception of an additional free parameter of scale height, while fitting the rotational velocity. We notice from the channel maps, two of which are shown in Figure 5.5, that our observed simulations (in blue) and the model (in red) tally quite well. Overall, the agreement between the results obtained with both codes is satisfactory, especially in the outer regions which are of greatest interest to us.

Note that, since we have focused on obtaining the rotation curves and pressure support corrections in the outer regions, we do not make strong claims about

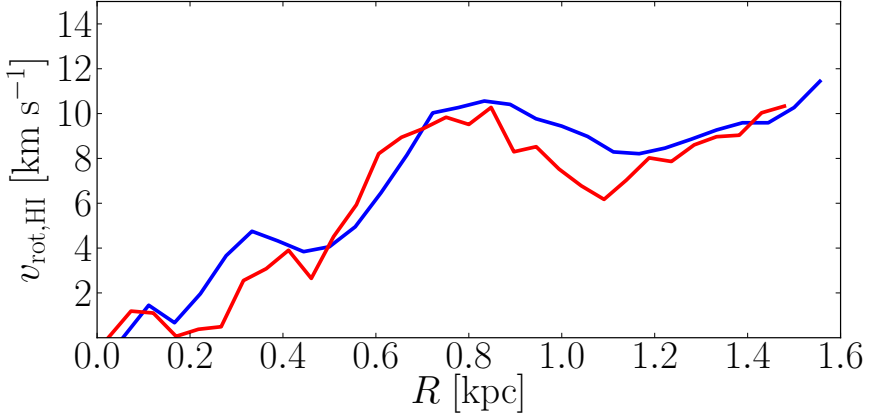


Figure 5.4: Rotation curves of one of the MoRiA dwarfs, obtained using GIPSY (in blue) and using ^{3D}Barolo (in red). The simulation is the one shown in Figure 5.3a.

the rotation in the central regions. To investigate i.e. the universality of dwarf galaxy rotation curves (Karukes & Salucci, 2017) or the radial acceleration relation (McGaugh *et al.*, 2016; Lelli *et al.*, 2017a) for the MoRiA galaxies, we would need to realistically obtain rotation curves at all radii. This lies beyond the scope of this paper.

5.3 Results

5.3.1 The $W_{50} - v_{\text{out,HI}}$ relation

An often wondered question is how W_{50} , which is relatively easy to measure, relates to the harder to obtain $v_{\text{out,HI}}$ (see e.g. Brook & Shankar, 2016; Ponomareva *et al.*, 2016). Figure 5.6 shows the relation between the two quantities for observed low-mass galaxies (compiled in P16). For galaxies with $v_{\text{out,HI}} \lesssim 70 \text{ km s}^{-1}$, the scatter on the relation becomes significant, and it typically holds that $W_{50}/2 < v_{\text{out,HI}}$. The MoRiA galaxies follow the trend of the observations, although some seem to be on the low-end of the data.

5.3.2 The $W_{50} - v_{h,\text{max}}$ relation

In their study of the TBTF problem in field dwarfs, P16 derived the average relation between the observed H I velocity width of galaxies, W_{50} , and the max-

5.3 Results

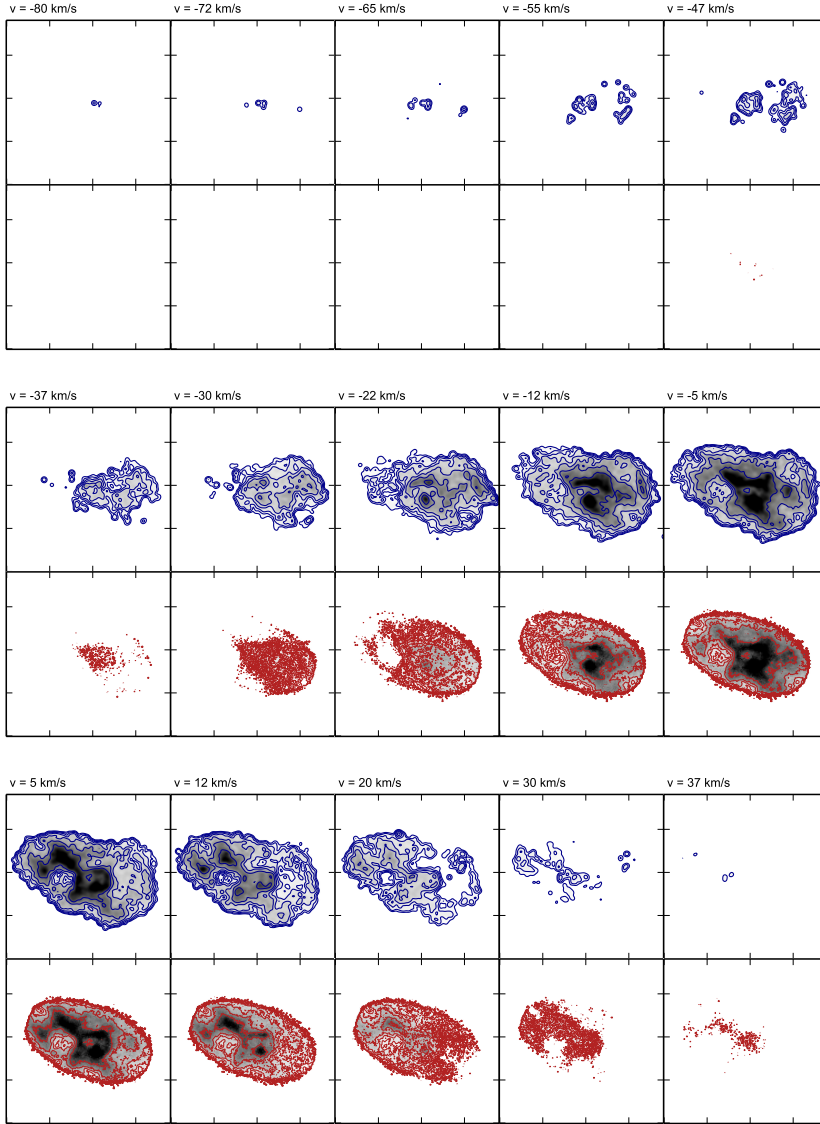


Figure 5.5: Channel maps of one of the MORIA dwarfs, in blue, and the model fit with 3^{D} Barolo, in red. The 3^{D} Barolo model reproduces the most salient features of the input data cubes. The simulation is the one shown in Figure 5.3a.

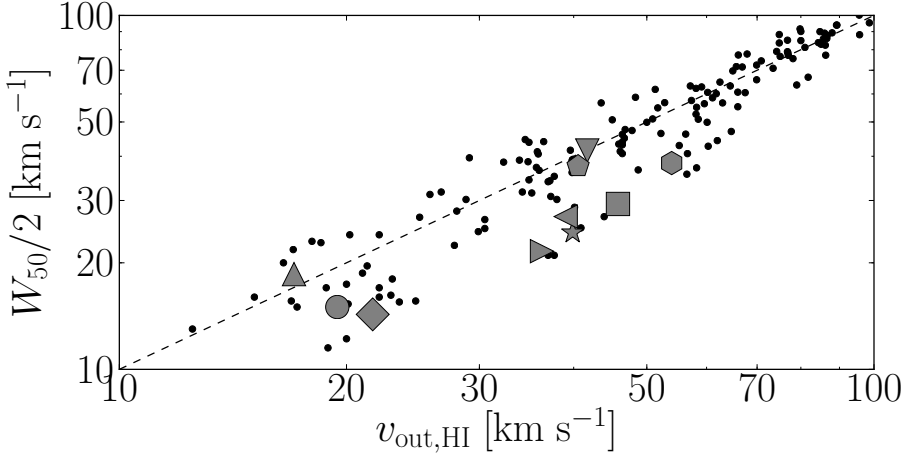


Figure 5.6: $W_{50}/2$ versus $v_{\text{out,HI}}$ for the MORIA dwarfs. Symbols are as indicated in Table 5.1. Observations, compiled in P16, are shown in black. The dashed line shows the case for both quantities being equal.

imum circular velocity of their host halos, $v_{h,\text{max}}^{\text{true}}$, such that the observed VF of galaxies (Haynes *et al.*, 2011; Klypin *et al.*, 2015, P16) can be reproduced within the Λ CDM cosmological model. The observed rotation velocity we use here is $W_{50}/2$.

Figure 5.7 shows the location of the MORIA dwarfs in the $W_{50} - v_{h,\text{max}}^{\text{true}}$ plane. The MORIA dwarfs follow the average relation derived in P16 very well, a fact that ensures that MORIA dwarfs are produced at the correct number densities as a function of their W_{50} (i.e., the MORIA simulation reproduces the observational VF). Similar results were also obtained by Macciò *et al.* (2016) based on the NIHAO hydrodynamical simulations (Wang *et al.*, 2015) and by Brooks *et al.* (2017) based on a set of hydrodynamic simulations of galaxy formation carried out by Governato *et al.* (2012), Brooks & Zolotov (2014) and Christensen *et al.* (2014).

However, reproducing the observational VF alone does not necessarily mean that the cosmological problems faced by Λ CDM on small scales have been resolved. In particular, a successful simulation must also be able to reproduce the internal, spatially resolved, kinematics of observed dwarfs. This is a crucial point, since the inconsistency between the predicted velocity profiles of simulations that are able to reproduce the observational VF, and the measured outermost-point rotational velocities of small dwarfs is at the heart of the TBTF problem (e.g. Papastergis *et al.* 2016, see also Trujillo-Gomez *et al.* 2016; Schneider *et al.* 2016).

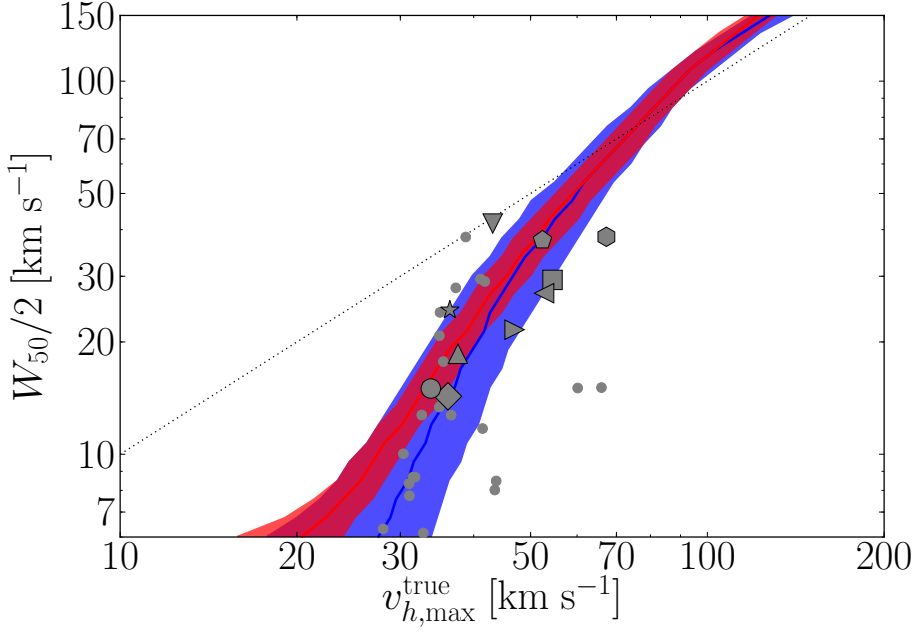


Figure 5.7: The MoRiA dwarfs in the $W_{50} - v_{h,\max}$ plane. Gray symbols are the simulations for which we resolved rotation curves are available, with their symbols as indicated in Table 5.1. Gray dots indicate MoRiA simulations not explicitly discussed. The red and blue lines are the P16-relations derived from different observational datasets, with the bands around them representing their uncertainty.

5.3.3 Halo profile fitting

We use the NFW profile and the DC14 profile, which is given by the expression (Di Cintio *et al.*, 2014a)

$$\rho_{\text{DC14}}(R) = \frac{\rho_s}{\left(\frac{R}{R_s}\right)^\gamma \left(1 + \left(\frac{R}{R_s}\right)^\alpha\right)^{(\beta-\gamma)/\alpha}}, \quad (5.7)$$

where R_s is scale length and ρ_s is a multiple of the density at radius $R = R_s$. While the NFW profile was derived from dark-matter only simulations, the DC14 profile takes the halo response to baryonic effects into account. The α , β , and γ parameters are set by the star formation efficiency of the galaxy (quantified by the ratio of stellar to halo mass, $M_\star/M_h$). For $\alpha = 1$, $\beta = 3$, and $\gamma = 1$, the DC14 profile coincides with the NFW profile. If the stellar mass of the galaxy is known, both profiles have only two free parameters: the halo mass M_h and the halo concentration $c = R_{\text{vir}}/R_s$, with R_{vir} the virial radius.

P16 fitted both these profiles to the velocity measured at the outermost H I point of each galaxy (data taken from Begum *et al.* (2008a); de Blok *et al.* (2008); Oh *et al.* (2011); Swaters *et al.* (2009, 2011); Trachternach *et al.* (2009); Kirby *et al.* (2012); Côté *et al.* (2000); Verheijen & Sancisi (2001); Sanders (1996); Hunter *et al.* (2012); Oh *et al.* (2015); Cannon *et al.* (2011); Giovanelli *et al.* (2013); Bernstein-Cooper *et al.* (2014)). They fixed the halo concentration to the mean cosmic value ($\log_{10} c = 0.905 - 0.101 \log_{10}(M_h/(10^{12} h^{-1} \text{M}_\odot))$; Dutton & Macciò, 2014), leaving only the halo mass as a free parameter. From the fitted profile, they compute the maximum circular velocity of each galaxy’s host halo. In a successful cosmological model, individual galaxies should have $W_{50} - v_{h,\text{max}}^{\text{fit}}$ data-points that agree with the average $W_{50} - v_{h,\text{max}}^{\text{true}}$ relation that is needed to reproduce the observed VF (blue and red bands in Figure 5.7). As shown by P16, all is not well: a sizable fraction of low-mass galaxies fall to the left of the expected $W_{50} - v_{h,\text{max}}^{\text{true}}$ relation. In other words: the halo circular velocity implied by their H I kinematics is too low.

We exactly replicate this analysis for the MORIA dwarfs and show the results in Figure 5.8. Although there are fewer MORIA dwarfs than in the P16 sample, the result is broadly the same: the $W_{50} - v_{h,\text{max}}^{\text{fit}}$ relation is inconsistent with the expected $W_{50} - v_{h,\text{max}}^{\text{true}}$ relation. The simulations with the highest $v_{h,\text{max}}^{\text{fit}}$ -values seem to lie on the low end of, or even slightly below, the data points. This can be attributed to their lower-than-average W_{50} -values (see Figure 5.6). Another explanation is their smaller-than-average R_{out} values (see Figure 5.2a), since for smaller radii, the uncertainties on $v_{\text{out,H I}}$ will be extrapolated to large uncertainties on $v_{h,\text{max}}^{\text{fit}}$. The two (low W_{50} and small R_{out} values) most likely work together and probably come hand-in-hand. Indeed, for smaller H I bodies, the potential will be traced at smaller radii, resulting in a lower W_{50} .

5.4 Discussion

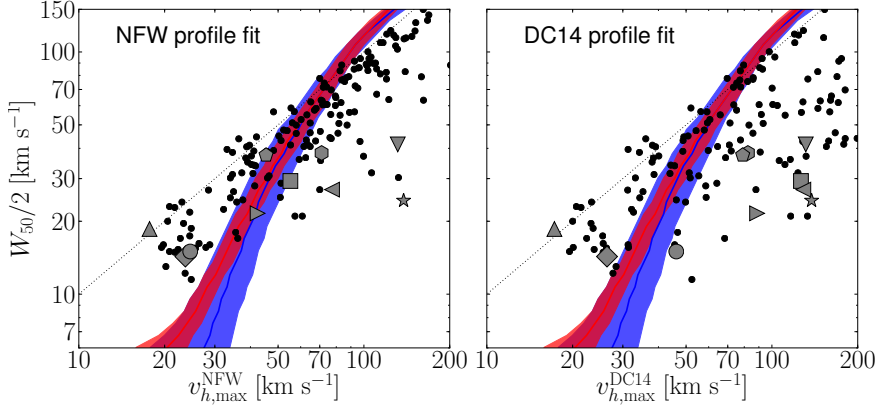


Figure 5.8: Results from fitting an NFW (top panel) and DC14 (bottom panel) to the outer-most point of the rotation curves of the MoRiA simulations using a fixed halo concentration. Symbols are as indicated in Table 5.1. This is compared to the results from P16 (in black). Red and blue lines and bands are the same as in Figure 5.7.

Analyzed in this way, one would be driven to the conclusion that the MoRiA dwarfs do not follow the $W_{50} - v_{h,\max}^{\text{true}}$ relation required for the Λ CDM halo VF to match the observed galactic VF and, therefore, that they suffer from the TBTF problem. The crucial difference between Figure 5.7 and Figure 5.8 is the fact that in the former Figure, the maximum halo velocity ($v_{h,\max}^{\text{true}}$) is computed directly from the enclosed mass profile of each simulated galaxy, while in the latter Figure, $v_{h,\max}^{\text{fit}}$ is computed by fitting the mock H I kinematics of each simulated galaxy.

5.4 Discussion

5.4.1 Concentration fitting

To be able to fit the two-parameter density profiles given by equations (1.5) and (5.7) to only two points (the central and outermost measured point of the rotation curve), P16 kept the concentration c of the halos fixed at the mean cosmic value and used the halo mass M_h as a free parameter. Another choice would be to fix M_h using the stellar mass and an abundance matching relation and to keep the concentration c as a free parameter.

In Figure 5.9, we show the $W_{50} - v_{h,\max}^{\text{fit}}$ relation obtained by fitting NFW and DC14 profiles to the P16 dataset, using the concentration c as a free parameter

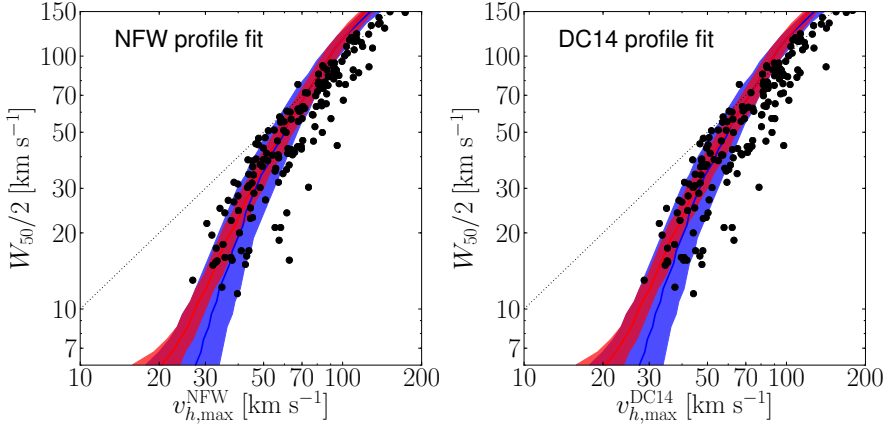


Figure 5.9: Results from fitting an NFW (top panel) and DC14 (bottom panel) to the outer-most point of the rotation curves of the observations used in Papastergis & Shankar (2016) using a fixed halo mass, keeping the halo concentration as a free parameter. The halo mass is calculated from their stellar mass using the abundance matching relation of Moster *et al.* (2013). Red and blue lines and bands are the same as in Figure 5.7.

with the halo mass M_h set by the stellar mass and the Moster *et al.* (2013) abundance matching relation. This way, the observed galaxies adhere much more closely to the expected $W_{50} - v_{h,\max}^{\text{true}}$ relation. NFW and DC14 profiles now actually produce very similar results.

In Figure 5.10, we compare the concentrations of the P16 galaxies retrieved in this way with the mass-dependent cosmic mean value derived from cosmological simulations (Dutton & Macciò, 2014). The frequency distribution of the concentration values is well approximated with a log-normal distribution function. Both for an NFW and a DC14 fit, the scatter is $\sigma \approx 0.25 - 0.3$ dex. This is significantly larger than the scatter on $\log(c/\langle c \rangle)$ found in cosmological simulations, where $\sigma \sim 0.13$ (Dutton & Macciò, 2014).

In Figure 5.10, we distinguish between galaxies with $W_{50}/2 < 30 \text{ km s}^{-1}$ (red data-points) and $W_{50}/2 > 30 \text{ km s}^{-1}$ (blue data-points), for the fits with the NFW and DC14 profiles. We choose the 30 km s^{-1} split because this is the rotation velocity below which the TBTF problem becomes apparent. Clearly, the high- W_{50} galaxies have higher concentrations than expected while the low- W_{50} dwarfs have lower concentrations, with a hint of an anticorrelation between M_h and concentration for the low W_{50} galaxies. This is to be expected: higher-mass galaxies must have lower concentrations in order to have low circular velocities.

5.4 Discussion

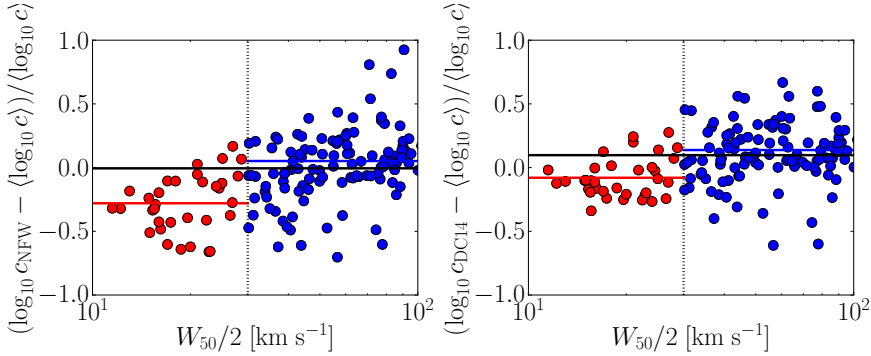


Figure 5.10: Logarithmic difference between the fitted concentration and the one expected from cosmological dark-matter only simulations (Dutton & Macciò, 2014) of the P16 sample of galaxies. Top panel: concentration obtained by fitting the NFW density profile to the kinematic data; bottom panel: concentration obtained using the DC14 density profile. Red and blue symbols indicate galaxies with low and high H I rotation velocities, respectively. The red and blue lines indicate the mean concentration of both subsamples; the green line indicates the mean concentration of the full sample.

To see whether the average of each subsample differs significantly from the cosmic mean value, we ran a t -test on both populations. We find p -values of 0.025 and 4.2×10^{-8} for the galaxies with respectively high and low W_{50} . The full sample has a mean concentration consistent with the Dutton & Macciò (2014) simulations, unlike the subsamples separately. For the DC14 profile, the same trend is found, with all the averages slightly higher than for the NFW profile. The p -values for the t -test are 2.0×10^{-7} for the entire sample and 1.3×10^{-10} and 0.0047 for the high and low circular velocity samples respectively. Employing the DC14 density profile yields concentration estimates that are inconsistent with the Dutton & Macciò (2014) simulations, both for the full sample and the subsamples.

Both the large scatter and the offsets are probably due, at least in part, to uncertainties on the (extrapolated) low-mass end of the $M_\star - M_h$ relation that was used to derive the halo mass from the stellar mass. In the mass regime we are interested in, the scatter on the $M_\star - M_h$ relation is expected to be substantial (e.g. Sales *et al.*, 2017) and, using our approach, this translates in an increased scatter on the concentration parameter. Moreover, there is great variation among the different published $M_\star - M_h$ relations in the regime of dwarf galaxies ($M_h \approx 10^{10} M_\odot$). Simulations also show a large scatter in stellar mass for these type of halos (e.g. Figure 7 in V15). We redid our analysis adopting different $M_\star - M_h$ relations: using the relation of Guo *et al.* (2010), we reach

the same conclusions as for the relation of Moster *et al.* (2013); when using the relation of Behroozi *et al.* (2013), the fitted concentrations are more in line with the predictions from Λ CDM, however they do not follow the P16-relation.

Katz *et al.* (2017) fitted an NFW and a DC14-profile to 147 SPARC-galaxies, taken from a sample of 175 galaxies with extended H I rotation curves (Lelli *et al.*, 2016a). They conclude that the fitted halo masses and concentrations for the DC14-profile are in line with the predictions from Λ CDM. The difference with our results is that their conclusions is that they only discuss their entire sample, which consists mostly of high-mass galaxies. They also use full rotation curves to fit the halo profile to the galaxies, allowing them to fit the halo mass and concentration simultaneously.

By fitting a *coreNFW* profile (Read *et al.*, 2016) to full rotation curves of a subset of the Little THINGS galaxies (Iorio *et al.*, 2017), Read *et al.* (2017) find that these isolated dwarf galaxies inhabit halos consistent with the abundance matching relation of Behroozi *et al.* (2013) and, as such, do not find a TBTF for isolated galaxies *at all*. These conclusions would change when assuming a different $M_\star - M_{\text{halo}}$ relation, as they remark in their Appendix C. Even so, we still find that using the relation of Behroozi *et al.* (2013), the observations do not follow the P16-relation.

5.4.2 Does the H I rotation curve trace the potential?

The fact that $v_{h,\text{max}}^{\text{true}}$ and $v_{h,\text{max}}^{\text{fit}}$ have different values, might be explained by an observational effect: the H I rotation curve of dwarf galaxies does not exactly follow the underlying potential.

Indeed, as one can see in Figure 5.11, the H I circular velocity profiles ($v_{\text{circ}}^{\text{obs}}(R)$) of the MORIA dwarfs are quite different, even after correcting for pressure support, from the true circular velocity profiles ($v_h^{\text{true}}(R)$). More often than not, *the outer H I rotation velocity data-point falls significantly below the true value of the local circular velocity*. It is important to keep in mind that the preceding statement is not directly related to the fact that rotational velocities derived from the linewidth of the HI profile of dwarf galaxies, $W_{50}/2$, underestimate the maximum circular velocity of the host halo, $v_{h,\text{max}}^{\text{true}}$, a result that has already been reported by Macciò *et al.* (2016) and Brooks *et al.* (2017). In fact, the linewidth derived H I velocity probes radii much smaller than the radius where the host halo rotation curves peaks, and thus there is no guarantee that the two quantities should be the same. On top of this, it is different from the fact that the H I rotation curve is still rising at its outer-most radius and thus does not trace $v_{h,\text{max}}^{\text{true}}$ (Brook & Di Cintio, 2015b; Ponomareva *et al.*, 2016). What we demonstrate here instead is that the circular velocity computed from spatially resolved H I data underestimates the true circular velocity *at the same radius*.

5.4 Discussion

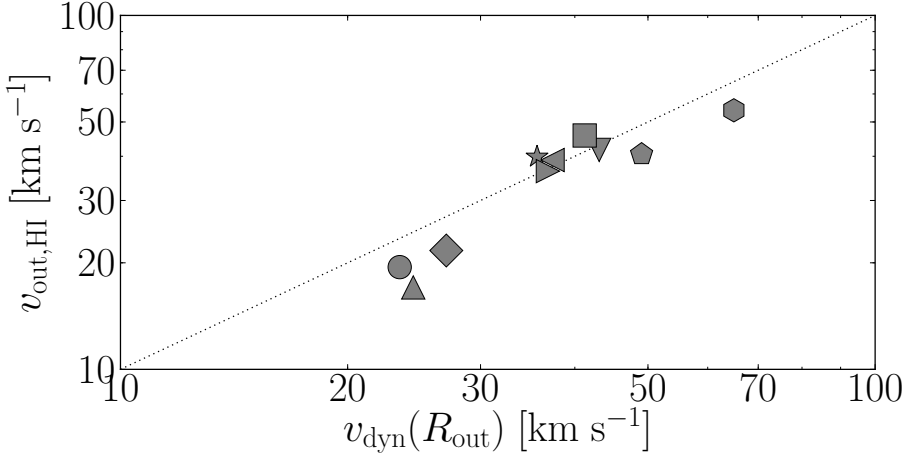


Figure 5.11: The measured H I rotation at R_{out} compared to the dynamical circular velocity, as expected from the enclosed mass. Symbols are as indicated in Table 5.1.

Of course, there are only ten MORIA dwarf galaxies with resolved rotation curves and a bigger sample of simulated dwarfs is definitely needed to fully explore this issue. But if this explanation holds water, it would explain why the halo fitting using a fixed concentration fails: we are not fitting to the actual halo velocity at this radius.

In the previous Section 5.4.1, we redid the halo fitting, but now fixing the halo mass using an abundance matching relation and keeping the halo concentration as a free parameter. The resulting concentrations do not seem to be drawn from the distribution predicted by Λ CDM, especially for galaxies with low W_{50} . If the observed H I rotation curve does not trace the potential, this would explain the seemingly wrong population of concentrations.

Valenzuela *et al.* (2007) and Pineda *et al.* (2017) have also applied a tilted-ring method to derive the H I rotation curve, to investigate whether dwarf galaxies have dark matter cusps or cores. They studied galaxies with an idealized set-up and both find that the observed H I rotation underestimates the gravitational potential, however only in the inner regions. Here, using more realistic dwarf galaxies, we show that the idea that the H I rotation is not necessarily a good tracer for the underlying gravitational potential of dwarf galaxies is not necessarily confined to the inner regions of galaxies, but extends over their entire body. One crucial question here is what causes this substantial underestimate of the local circular velocity in observational measurements of the HI kinematics. We

attribute this to the fact that the assumptions underlying the tilted-ring fitting method and the correction for pressure support are not met in the case of low-mass dwarf galaxies: their atomic ISM simply does not form a relatively flat, dynamically cold disc. Rather, they have a vertically thick ($\langle q \rangle \sim 0.5$), dynamically hot, continuously stirred atomic ISM with significant substructures that is not in dynamical equilibrium in the gravitational potential. The detailed analysis of the vertical structure of the HI disks of MORIA dwarfs and of non-circular motions in their velocity fields will be the focus of a separate publication (Verbeke et al. in prep.)

5.4.3 Stellar kinematics

In accordance with the analysis of P16, we have used H I ordered motions to get an idea of the underlying gravitational potential. Late-type dwarf galaxies are typically dispersion-supported (Kirby *et al.*, 2014; Wheeler *et al.*, 2017). So the stellar velocity dispersion σ_* of a late-type dwarf can also be used to obtain a mass estimate, by using

$$M_{\sigma_*}(R) = 3\sigma_*^2 R G^{-1}. \quad (5.8)$$

For our simulations, we want to calculate σ_* in the same way as is done observationally (Kirby *et al.*, 2014). In the same vein as Vandenbroucke *et al.* (2016), we weigh the average with the number of red giant branch (RGB) stars expected in each stellar particle (using the stellar evolution models of Bertelli *et al.*, 2008, 2009). As shown in Figure 5.12a, the enclosed mass within R_{out} inferred from the stellar velocity dispersion agrees reasonably well with the one inferred from the H I kinematics ($M_{\text{circ,HI}} = v_{\text{out,HI}}^2 R_{\text{out}} G^{-1}$). The mass estimated from stellar velocity dispersions agrees typically within $\sim 30\%$ with the mass inferred from the HI rotation curve. Two simulations however have a relative difference of $\sim 50\%$. Figure 5.12b shows the measured stellar velocity dispersion within R_{out} in function of the dynamical one, σ_{dyn} , as expected from the enclosed mass. Note that we have only included simulations with at least 100 stellar particles within R_{out} , to get a good measure of σ_* . Contrary to Figure 5.11, the lowest-mass MORIA dwarfs do not systematically have lower observed stellar velocity dispersions than dynamical ones. However, over our entire sample, the majority of the simulated dwarfs have $\sigma_* < \sigma_{\text{dyn}}$. This implies that in most cases, their dynamical mass would be underestimated from observed stellar kinematics.

Given this, the results presented in this paper might also be extended to the Too-Big-To-Fail problem for satellite galaxies. However, other effects play an important role. The presence of H I and active star formation (and thus stellar feedback) in field dwarfs will influence the stellar kinematics through dynamical heating or cooling. Satellites are devoid of H I but will on the other hand be influenced by the tidal field from their host galaxy. A similar study of simulated

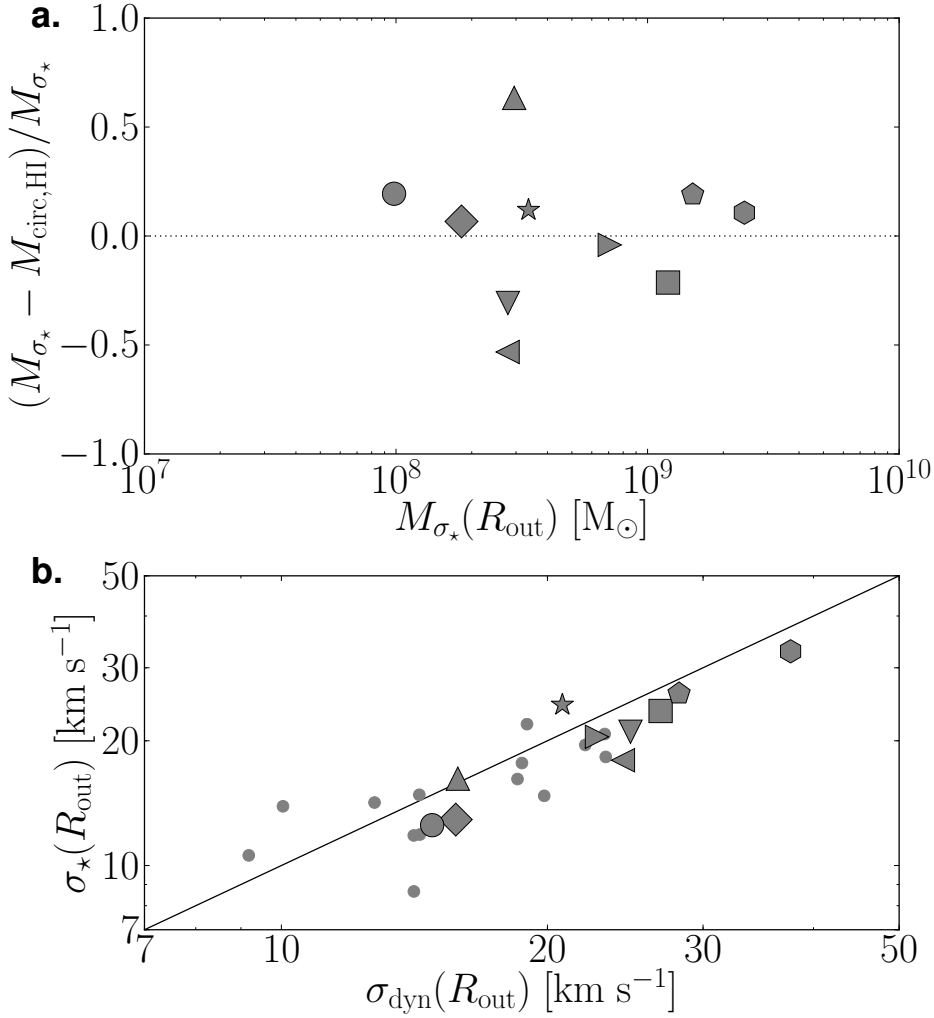


Figure 5.12: **a.** The relative difference between the enclosed mass within R_{out} inferred from stellar velocity dispersions σ_* and H I circular motions. **b.** The stellar velocity dispersion within R_{out} compared to the dynamical velocity dispersion, as expected from the enclosed mass. Symbols are as indicated in Table 5.1, with dots the MORIA galaxies without resolved rotation curves.

satellite galaxies is thus necessary to see if their stellar kinematics underestimate the halo mass.

5.4.4 Are disturbed velocity fields realistic?

As can be seen from Figure 5.3, the H I distributions of the MORIA galaxies are generally quite disturbed. This is also the case for real dwarf galaxies with velocity widths $W_{50}/2 \lesssim 30 \text{ km s}^{-1}$; see e.g. Leo P (Bernstein-Cooper *et al.*, 2014), CVndwA, DDO 210, and DDO 216 (Oh *et al.*, 2015). However, galaxies with larger velocity widths, $W_{50}/2 \sim 40 - 70 \text{ km s}^{-1}$, typically display regular velocity fields and low H I velocity dispersions, $\sigma_{\text{HI}} \lesssim 12 \text{ km s}^{-1}$ (e.g. Kirby *et al.*, 2012; Iorio *et al.*, 2017). In contrast, the most massive MORIA dwarf that we have analyzed, M-10, has a fairly disturbed velocity field and a relatively large velocity dispersion, $\sigma_{\text{HI}} \gtrsim 20 \text{ km s}^{-1}$. Even though we cannot draw reliable conclusions from this one object alone, it is possible that this indicates that the efficiency of stellar feedback in the MORIA simulation is too strong. Note that most state-of-the-art hydrodynamic simulations of dwarf galaxy formation (e.g. Hopkins *et al.*, 2014; Wang *et al.*, 2015; Sawala *et al.*, 2016a) have more efficient feedback schemes than MORIA, so this could represent a general issue for (dwarf) galaxy simulations. At the same time however, the MORIA simulation reproduces successfully the sizes and thickness of H I disks of observed dwarfs (Figures 5.1 and 5.2). Moreover, in Figure 9 of V15 we have shown that the spatial distribution of H I in MORIA dwarfs has similar power spectrum slopes as those measured for LITTLE THINGS galaxies (Zhang *et al.*, 2012; Hunter *et al.*, 2012). We leave the investigation of this, including the effect of e.g. beam size, for future research. In any case, this does not change our conclusions in any way since these are based on the galaxies with $W_{50}/2 \lesssim 30 \text{ km s}^{-1}$.

5.5 Conclusions

We have used the MORIA simulations of dwarf galaxies with realistic H I distributions and kinematics to investigate the Too Big To Fail problem for late-type field dwarfs.

We showed that the MORIA dwarfs follow the relation between H I line-width and halo circular velocity, derived by Papastergis & Shankar (2016), that is required for the Λ CDM halo velocity function to correspond to the observed field galaxy width function. This means that, given the number density of halos formed in a Λ CDM universe, the MORIA simulations reproduce the observed galactic velocity function. In other words: there is no missing dwarfs in the MORIA simulations.

We then constructed resolved H I rotation curves, including corrections for pressure support, for ten of the MORIA dwarf galaxies. We used our mock H I rotation

5.5 Conclusions

curves to replicate the analysis of Papastergis & Shankar (2016) and fitted NFW and DC14 density profiles (with fixed concentration parameter) to the outermost point of these measured rotation curves to derive an observational estimate for the maximum halo circular velocity of each MORIA galaxy. Using this estimate for the circular velocity, the MORIA dwarfs, like the real dwarf galaxies analyzed by Papastergis & Shankar (2016), fail to adhere to the relation between H I line-width and halo circular velocity that is required for the Λ CDM halo velocity function to correspond to the observed field galaxy width function. In other words: using only quantities derived from observations, dwarf galaxies (both real and simulated) experience the Too Big To Fail problem.

What causes this difference between the results from fitting a halo profile to the outer-most point of the rotation curve and using the actual $v_{h,\max}^{\text{true}}$ -value derived directly from the mass distribution?

Comparing the H I rotation curves of the MORIA dwarf galaxies with their theoretical halo circular velocity curves, we see that they can differ significantly. The circular velocities derived from the H I kinematics of MORIA dwarfs with H I rotation velocities below $\sim 30 \text{ km s}^{-1}$ are typically too low. This results in too low a $v_{h,\max}^{\text{fit}}$ value at a fixed concentration c . The TBTF problem thus results, at least partially, from the fact that for galaxies in this regime, its halo mass can not readily be inferred from its (H I) kinematics. Indeed, based on their kinematics, galaxies with $W_{50}/2 \lesssim 30 \text{ km s}^{-1}$ would seem to inhabit less massive halos than they actually do. However, under the assumptions of Λ CDM, the P16-relation does provide an estimate of the true halo circular velocity $v_{h,\max}^{\text{true}}$ in function of a galaxies H I linewidth W_{50} .

We attribute this effect to that fact that the atomic interstellar medium of low-mass dwarfs simply does not form a relatively flat, dynamically cold disc whose kinematics directly traces the underlying gravitational force field. Another explanation might be that the H I velocity fields are too irregular to infer its halo mass from its kinematics. This is true for most of the simulated galaxies presented in this work, as well as for observed low-mass galaxies.

The stellar feedback efficiency will influence both these H I-thickness of the galaxies as well as how messy the velocity fields are. Thus, how much energy is actually injected in the ISM by stellar feedback is an important issue in the discussion of the Too Big To Fail problem.

Ongoing & future work

The pursuit of knowledge is
hopeless and eternal... HOORAY!

Professor Hubert J. Farnsworth

HERE we give a short overview of projects we are currently working on and that have thus not been published yet. We end with an outlook to future research.

6.1 Colour-Magnitude Diagrams

To further bridge the gap between simulations and observations, we would also like to adjust the way SFHs are derived from simulations. Now, we just use the particle data present in our simulated data: each stellar particle has an age, so the SFH is easily derived. Observationally, it is generally not possible to assess the age of an individual star. And even if it was, deriving it for the millions or billions of stars in each galaxy would not be feasible. However, stars are not born individually. A stellar population born at the same time but with different masses will have a specific spectrum depending on its age and metallicity.

If the individual stars of a galaxy can be resolved, these can be presented in a colour-magnitude diagram (CMD). A single stellar population would occupy a specific region in a CMD, called an isochrone. As such, it is possible to derive the SFH of a galaxy from its CMD. There are some caveats, since the spectrum of a stellar population depends degenerately on its age and metallicity. Moreover, the theoretical stellar models are not as well determined as one would like.

Using the BaSTI models (Pietrinferni *et al.*, 2013, Bag of Stellar Tracks and Isochrones), we can construct a CMD for our simulations. The models cover a broad range in metallicity ($10^{-5} \leq Z \leq 0.05$) with an α -dependence (both scaled-solar and $[\alpha/\text{Fe}] = 0.4$ models exist). We chose the values with the mass loss

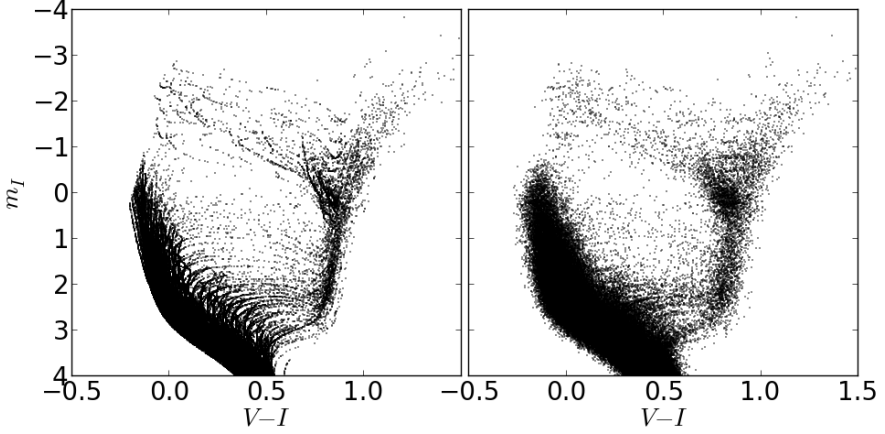


Figure 6.1: Example colour magnitude diagram for one of the MORIA simulations. (a) Without convolving, (b) with convolving with the errors in the V - and I -band following a Gaussian distribution with $\sigma = 0.05$.

parameter $\eta = 0.4$. Each stellar particle in our simulations represents a simple stellar population with a certain age, metallicity and $[\alpha/\text{Fe}]$, and we assume the isochrone with the values closest to these.

We then need to sample the isochrone associated with each stellar particle with individual stars. This can be done by inverting the cumulative version of the Chabrier IMF (equation 2.23):

$$M(r) = 10^{(\text{erf}^{-1}(Nr/D + \text{erf}(A(B-1)))/A - B)} \quad \text{if } r \leq \Psi(1) \quad (6.1)$$

$$= (1 - 1.3N(r - \Psi(1))^{-1/1.3}) \quad \text{if } r \geq \Psi(1), \quad (6.2)$$

where A and B are the same as in equation 2.23), N is a normalization factor, $D := \frac{C\sqrt{\pi}}{2B\log_{10}(e)}$, $\Phi(m)$ is the cumulative IMF, erf is the error function, and r is a parameter between 0 and 1. If we then randomly draw values for $r \in [0, 1)$ and associate it with a mass using equation 6.2, this will follow the IMF (equation 2.23).

With each sample star we can associate V -band and $V - I$ colour by linearly interpolating on the values tabulated in the isochrone.

The stellar particles have a certain initial mass, thus without taking mass loss into account, so we can sample individual stars until this total mass is reached.

6.1 Colour-Magnitude Diagrams

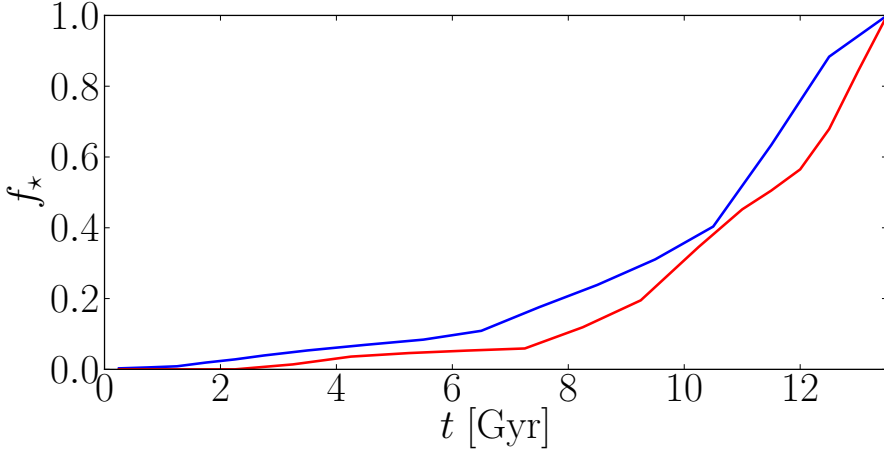


Figure 6.2: SFHs of DGe15e9, derived from the simulation data (red) and from the CMD (blue).

Not every sampled star will contribute to the CMD, since the IMF does not take the finite lifetime of stars into account. Massive stars can thus be sampled and will contribute to the total sampled mass, but will not be put on the CMD. The technique for populating the IMF will typically lead to an oversampling, since we can not make sure that the total mass is sampled exactly. To prevent this error to grow too large while sampling all the stellar particles, we keep track of the error after a particle has been fully sampled and add this to the mass that has already been sampled when starting the process for the next particle. This process ensures that the overshoot for the entire simulation will be smaller than $70 M_{\odot}$, the upper limit for our IMF. Furthermore, we can convolve the obtained theoretical values to account for uncertainties and observational errors. An example of a colour magnitude diagram obtained using this method for one of the MORIA simulations can be seen in Figure 6.1.

To derive the SFH from a CMD, we use the same method as described in Bernard *et al.* (2012). We compare the SFH obtained directly from the data to one inferred from the CMD for one of the MORIA simulations (DG15e9) in Figure 6.2. As can be seen, the comparison is quite good. DG15e9 had assembled most of its stellar mass very recently. Since young stellar populations are better constrained than older ones, it would be interesting to see how these results differ between simulated galaxies with an earlier mass assembly.

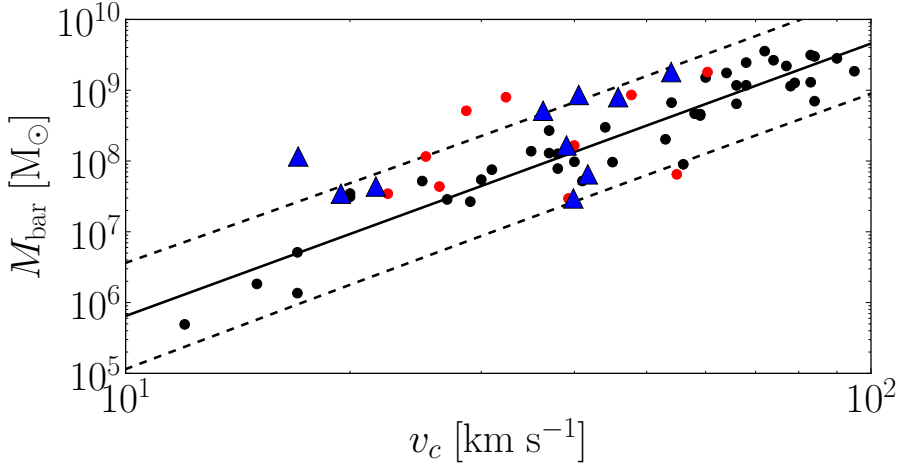


Figure 6.3: Baryonic Tully-Fisher relation for 10 MORIA galaxies using resolved H I rotation curves (in blue) compared to the same galaxies using $v_c = W_{20}$ (in red). Black dots are observations.

6.2 Baryonic Tully-Fisher Relation revisited

In Section 4.3.1, we discussed how the MORIA simulations compare to observations in terms of the BTFR. As seen in Figure 4.2, they follow the observed relation, but they do seem to have larger than observed scatter. The circular velocity plotted for the MORIA dwarfs was derived from their H I spectrum ($v_c = W_{20}/2$). However, the observational values are taken from resolved H I rotation curves, typically at the radius where the rotation curves becomes flat (V_f ; McGaugh, 2012). Figure 6.3 shows the BTFR for MORIA galaxies for which we have obtained resolved rotation curves. At first glance, it seems that the scatter is slightly decreased: now only one galaxy falls outside the 3σ confidence interval instead of two. However, more than 10 simulated galaxies would need to be examined to make strong claims about this. Furthermore, since we have obtained all quantities involved in the BTFR in an observationally motivated way, it should be possible to add “observational” errors to the simulations as well. Observationally, the errors are consistent with a very small intrinsic scatter on the BTFR (Lelli *et al.*, 2016b), also when using linewidth data (Papastergis *et al.*, 2016). It remains to be seen if this can be reproduced within simulations.

6.3 Star formation efficiency

As can be seen from Figure 4.6, for a fixed halo mass the MoRiA simulations show a large variation in star formation efficiency (SFE; $M_\star/M_{\text{halo}}$) for $M_{\text{halo}} \lesssim 10^{10} M_\odot$. To see what might cause these different SFEs, we have selected those MoRiA dwarfs with $M_{\text{halo}} \approx 10^{10} M_\odot$ and see how massive their dark matter halos are at different times. The results are shown for 1 Gyr, 2 Gyr, and 6 Gyr in Figure 6.4. For all times, there seems to be a trend that halos that formed earlier have higher SFEs. The correlation is strongest for 6 Gyr (especially if we consider the two galaxies with the highest SFE as outliers). To incorporate the effects of all different times, we also look at M_{avg} , which is the halo mass averaged out over different time bins (with increments of 50 Myr). M_{avg} correlates better with the SFE than the halo mass at any individual time. The SFE of a dark matter halo thus seems to be influenced by its halo assembly over cosmic time.

The correlation is not perfect, so there might be extra effects, other than the stochasticity of star formation, that play a role. These results also raise the question why this assembly history plays such an important role in low-mass galaxies and less in more massive ones, where the SFE of a halo is very well determined by its mass.

6.4 Halo modification

It was already shown in Cloet-Osselaer *et al.* (2012) that the inclusion of baryonic physics in our models has a drastic influence on the dark matter density in the inner regions. To study this, one can fit a power law

$$\rho \propto r^\alpha \quad (6.3)$$

to the DM density in the region $0.01 < R/R_{\text{vir}} < 0.02$, with R_{vir} the virial radius. It was found that α strongly depends on the star formation efficiency $M_\star/M_{\text{halo}}$. Based on the MaGICC simulations (Brook *et al.*, 2012; Stinson *et al.*, 2013), Di Cintio *et al.* (2014b) proposed a prescription of the form

$$\alpha(x) = n - \log_{10} \left[\left(\frac{x}{x_0} \right)^{-\beta} + \left(\frac{x}{x_0} \right)^\gamma \right], \quad (6.4)$$

where $x = M_\star/M_{\text{halo}}$, $n = 0.132$, $x_0 = 8.89 \times 10^{-3}$, $\beta = 0.593$ and $\gamma = 1.99$. For $\log_{10}(M_\star/M_{\text{halo}}) < -4$, the logarithmic density slope is fixed to $\alpha = -1$. Tollet *et al.* (2016) on the other hand favour

$$\alpha(x) = n - \log_{10} \left[n_1 \left(1 + \frac{x}{x_1} \right)^{-\beta} + \left(\frac{x}{x_0} \right)^\gamma \right], \quad (6.5)$$

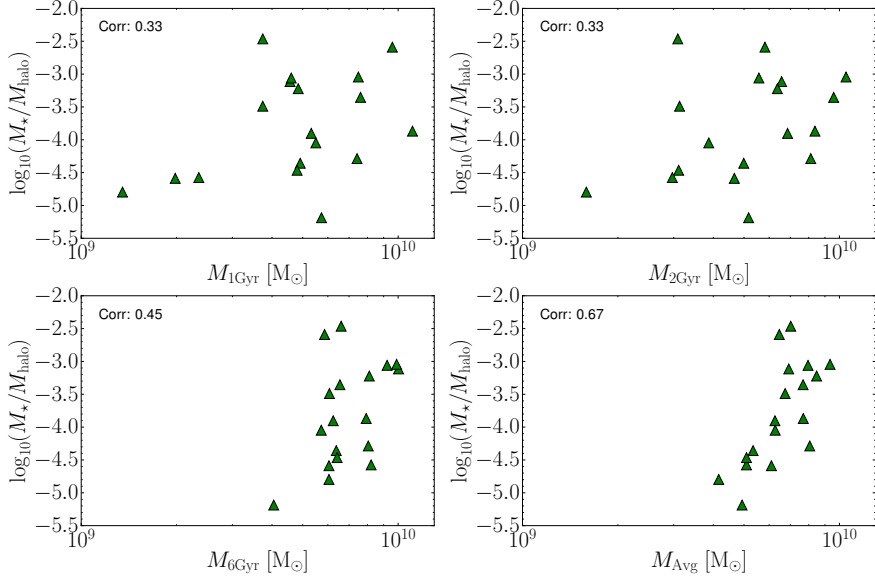


Figure 6.4: Star formation efficiency of the MORIA simulations with $M_{\text{halo}} \approx 10^{10} M_{\odot}$ in function of its dark matter halo mass at 1 Gyr (top left), 2 Gyr (top right), 6 Gyr (bottom left), and its halo mass averaged over cosmic time (M_{avg} , bottom right).

6.4 Halo modification

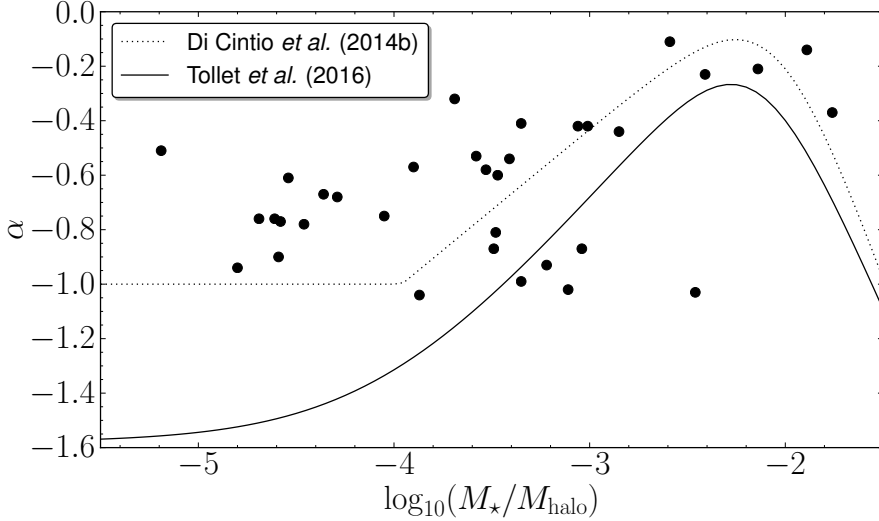


Figure 6.5: Inner density slopes of the dark matter halos of the MORIA galaxies compared to the relations of Di Cintio *et al.* (2014b) and Tollet *et al.* (2016).

$n = -0.158$, $n_1 = 26.49$, $x_0 = 8.77 \times 10^{-3}$, $x_1 = 9.44 \times 10^{-5}$, $\beta = 0.85$ and $\gamma = 1.66$. This has the advantage that it converges to $\alpha \rightarrow -1.58$ for $x \rightarrow 0$, which is close to the value expected from DMO simulations. Although the latter relation was based on the NIHAO simulations (Wang *et al.*, 2015), this is an updated version of the MaGICC project and thus share a very similar subgrid model. Chan *et al.* (2015) found the same overall trend using the FIRE simulations, but with some differences due to their different feedback implementation.

Figure 6.5 compares the α -values for the MORIA simulations to the above relations. For $M_*/M_{\text{halo}} \gtrsim 10^{-4}$, they roughly follow the proposed relations (albeit with one clear exception). Moreover, the peak of core creation at $M_*/M_{\text{halo}} \approx 10^{-2}$ is recovered as well. However, for $M_*/M_{\text{halo}} \lesssim 10^{-4}$, the MORIA galaxies have shallower density profiles than the MaGICC and NIHAO galaxies. The question is now if this has to do with the different initial conditions (cosmological zoom simulations versus our merger tree approach), with the specific implementation of the subgrid physics, or with something else. Since we include feedback from Pop III stars, we inject more energy from stellar feedback for a given stellar mass. The effects will be most drastic for galaxies with small stellar masses. Because of this, the shallower central density profiles for our galaxies with small stellar

masses are likely a physical result, rather than a numerical one.

6.5 Outlook

To end this thesis, we give a short outlook towards possible future research.

- In Chapter 5 we have studied the H I kinematics to obtain a measure for the circular velocity in the outer regions of the MORIA dwarfs. Obtaining rotation curves at all radii, would allow us to study the radial acceleration relation (McGaugh *et al.*, 2016; Lelli *et al.*, 2017a), the universality of rotation curves (Karukes & Salucci, 2017), and the cusp-versus-core problem (Moore, 1994; Flores & Primack, 1994).
- The role of gaseous infall to trigger starbursts in dwarf galaxies could be revisited using the more realistic MORIA simulations. This would allow for a proper comparison between real dwarfs, e.g. how their H I distribution and kinematics look like. Other trigger mechanisms, such as tidal interactions or (wet) mergers, could be explored as well.
- The analysis of the simulated galaxies can be made even more realistic. For example, we could use a radiative transfer code (e.g. SKIRT; Baes *et al.*, 2003; Camps & Baes, 2015) with the effects of metal line absorption and emission. The obtained spectrum can then be analyzed to get an observationally motivated measure of the metallicity of the galaxy.
- As a result of different people working on it over many years, our code could use some optimization. Both more well-designed algorithms as more efficient parallelization could drastically increase the efficiency of the code. In the coming years, computers will become increasingly more parallelized. An efficient simulation code is crucial to make optimal use of them.
- Closely related to the previous point, one might opt to start from a different N-body/hydro code than GADGET-2. A moving-mesh code such as SHADOWFAX (Vandenbroucke & De Rijcke, 2016) was shown to be more accurate than the SPH approach.
- It would be interesting to study the effects of the merger tree approach and solidify it as a well-motivated alternative to cosmological zoom simulations. To do this, a thorough comparison is needed between galaxies simulated with the merger tree approach and those formed in cosmological simulations is needed, using the same subgrid physics and resolution.

6.5 Outlook

- With the upcoming observational facilities such as the James Webb Space Telescope, the European Extremely Large Telescope, the Square Kilometer Array, and the Euclid mission, properties of galaxies and the large scale structure of the universe will be much more constrained. The merger tree approach can then be used to simulate individual galaxies with very high resolution. However, to study the large scale structure, cosmological simulation would need to be run. This further highlights the need for an efficient simulation code.

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