Exploring the nature of ISM turbulence in disc galaxies

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Abstract - English

Galaxy formation is a continuous process that started only a few hundred million years after the Big Bang. The first galaxies were very volatile, with bursts of star formation and disorganised gas motions. However, even as these galaxies evolved to have orderly rotating gas discs, the gas within the disc, referred to as the interstellar medium (ISM), still remained highly turbulent. In fact, the ISM is supersonically turbulent, meaning that the disorganised gas motion exceeds the speed of sound in the medium. This supersonic turbulence has been connected to several crucial properties related to galaxy evolution; for example, increasing (and decreasing in some regions) the ISM gas density, star formation, and gas mixing.

Many observations have shown that all of the gas phases in the ISM experience supersonic levels of turbulence, with line widths (an observational method to quantify the amount of turbulence) as high as \( \sigma_g \lesssim 100 \text{ km s}^{-1} \) in high-redshift (younger) disc galaxies, while local quiescent discs have \( \sigma_g \lesssim 40 \text{ km s}^{-1} \). However, the ISM contains a variety of gas phases that cover a wide range of temperatures and densities, which exhibit different levels of turbulence. For example, the warm ionised gas phase represents the upper limits quoted above, while colder denser gas only reaches \( \sigma_g \lesssim 40 \text{ km s}^{-1} \) and \( \sigma_g \lesssim 15 \text{ km s}^{-1} \) in high-redshift and local galaxies, respectively.

The physical processes driving this turbulence are not fully understood, but a combination of stellar feedback (e.g. supernova) and gravitational instability (e.g. during cloud collapse) have been suggested to provide a majority of the turbulent energy. In particular, stellar feedback is crucial in the formation of warm ionised gas and may therefore have a significant contribution on the turbulence within ionised gas. Furthermore, heterogeneous data of widely different galaxies (in terms of e.g. mass and size) at different resolutions (which causes artificial line broadening) complicates understanding the underlying cause.

A commonly used tracer of ionised gas is the H\( \alpha \) emission line and has been used extensively in high-redshift surveys. However, the contribution of the H\( \alpha \) signal comes from two primary sources: the radiatively ionised regions around massive newborn stars embedded in molecular gas (called H II regions) and diffuse ionised gas (DIG) filling the entire galactic disc. Observations have found that these two sources contribute, on average, roughly the same amount to the H\( \alpha \) signal (although with a large spread), but the levels of turbulence is starkly different; with the DIG being roughly 2-3 times more turbulent.
than the gas in HII regions.

Numerical simulations have come a long way and are now able to simulate entire disc
galaxies at parsec-scale resolution (in regions of interest). Furthermore, galaxy simulations
have been able to reproduce the level of turbulence observed in local and high-redshift
galaxies. Direct comparisons between numerical and observational studies are crucial to
understand the relevant physics driving observed correlations. However, numerical and
observational work have different data available and the reduction/analysis varies between
authors, and so diligence is required to perform qualitative comparisons.

In this work, I perform numerical simulations to investigate ISM turbulence in different
gas phases. My simulations model a Milky Way-like galaxy at two different redshifts
(using gas fraction as a proxy for redshift) and with/without stellar feedback physics, to
evaluate its impact. I perform mock observations to explore the relation between the star
formation rate and turbulence, and investigate what is driving this relation. Additionally, I
analyse the Hα emission line and compare the contribution in intensity and line broadening
(turbulence) from H II regions and DIG.
Abstract - Svenska

Galaxbildning är en kontinuerlig process som började bara några hundra miljoner år efter Big Bang. De första galaxerna var mycket volatila, med utbrott av stjärnbildning och oorganiserade gasrörelser. Men även efter att dessa galaxer utvecklade ordnade roterande gasskivor, förblev gasen inom skivan, kallat det interstellära mediet (ISM), fortfarande högt turbulent. Faktum är att ISM är supersoniskt turbulent, vilket innebär att de oorganiserade gasrörelserna överstiger ljudets hastighet i mediet. Denna supersoniska turbulens har kopplats till flera avgörande egenskaper relaterade till galaxutveckling; till exempel, öka (och i vissa regioner minska) ISM:ets gas densitet, stjärnbildning och gasblandning.

Många observationer har visat att alla gasfaser i ISM upplever supersoniska nivåer av turbulens, med linjebredder (en observationsmetod för att kvantifiera mängden turbulens) så höga som $\sigma_g \lesssim 100 \text{ km s}^{-1}$ i hög-rödförskjutnings (dvs. yngre) skivgalaxer, medan lokala lugna skivor har $\sigma_g \lesssim 40 \text{ km s}^{-1}$. Emellertid innehåller ISM olika gasfaser som täcker ett brett spektrum av temperaturer och densiteter, vilka uppnås olika nivåer av turbulens. Till exempel representerar den varma joniserade gasfasen de övre gränserna som nämns ovan, medan kallare, tätare gas endast når $\sigma_g \lesssim 40 \text{ km s}^{-1}$ och $\sigma_g \lesssim 15 \text{ km s}^{-1}$ i hög-rödförskjutnings och lokala galaxer, respektive.

De fysikaliska processer som driver denna turbulens är inte fullt förstådda, men en kombination av stellär feedback (t.ex. supernova) och gravitationsinstabilitet (t.ex. under molnkollaps) har föreslagits ge en majoritet av den turbulenta energin. I synnerhet är stellär feedback avgörande för bildandet av varm joniserad gas och kan därför ha ett betydande bidrag till turbulensen inom joniserad gas. Dessutom komplicerar heterogena data från mycket olika galaxer (i termer av t.ex. massa och storlek) vid olika upplösningar (vilket orsakar konstgjord linjebreddning) förståelsen av den underliggande orsaken.

En vanligt använd spårare av joniserad gas är H$\alpha$-emissionslinjen och har använts omfattande i undersökningar vid hög rödförskjutning. Emellertid kommer bidraget från H$\alpha$-signalen från två primära källor: de strålningsjoniserade regionerna runt massiva nyfödda stjärnor inbäddade i molekylär gas (kallade H$\text{II}$-regioner) och diffus joniserad gas (DIG) som fyller hela den galaktiska skivan. Observationer har funnit att dessa två källor bidrar, i genomsnitt, ungefär lika mycket till H$\alpha$-signalen (dock med en stor spridning), men nivåerna av turbulens är markant olika; med DIG ungefär 2-3 gånger mer turbulent än gasen i H$\text{II}$-regioner.
Numeriska simuleringar har kommit långt och kan nu simulera hela skivgalaxer med parsec-skala upplösning (i områden av intresse). Dessutom har galaxsimuleringar kunnat återskapa den nivå av turbulens som observerats i lokala och hög-rödförskjutningsgalaxer. Men numeriska och observationsbaserade arbeten har olika tillgängliga data och reduktion/analyser varierar mellan författare, och därför krävs noggrannhet för att göra kvalitativa jämförelser.

I detta arbete utför jag numeriska simuleringar för att undersöka ISM-turbulens i olika gasfaser. Mina simuleringar modellerar jag en Vintergatan-liknande galax vid två olika rödförskjutningar (användande gasfraktion som en proxy för rödförskjutning) och med/utan fysik för stellär feedback, för att utvärdera dess påverkan. Jag utforskar förhållandet mellan stjärnbildningshastigheten och turbulensen, och undersöker vad som driver detta förhållande. Dessutom analyserar jag Hα-emissionslinjen och jämför bidraget i intensitet och linjebreddning (turbulens) från HII-regioner och DIG.
Papers included and my contributions

Papers part of this thesis

Paper I:
"From giant clumps to clouds - III. The connection between star formation and turbulence in the ISM"
T. Ejdetjärn, O. Agertz, G. Östlin, F. Renaud, A. B. Romeo
Accepted to MNRAS in May 2022

Paper II:
"The origin of the Hα line profile in simulated disc galaxies"
T. Ejdetjärn, O. Agertz, G. Östlin, M. Rey, F. Renaud
Submitted to MNRAS in December 2023

My contributions

Paper I: The central idea of this project was an extension of my Master’s project. All numerical simulations of the galaxies were performed by me. I performed all of the analysis of the simulations (using self-written Python code), produced all of the figures, and wrote the majority of the text, with edits and feedback from my co-authors.

Paper II: The idea of this project was a follow-up of Paper I. The simulations were performed by OA, based on my simulations from Paper I. All of the analysis, figures, and text were produced by me, with edits and feedback from co-authors.
Papers not included

Here follows papers I have collaborated on during my work on this thesis, but that are not part of the main body of work.

Paper A:
"Ionizing radiation escape enabled by galaxy merger in reionization-era analog galaxy"
Acronyms

DM - Dark Matter
CDM - Cold Dark Matter
UV - Ultraviolet
IR - Infrared
ISM - Interstellar Medium
GMC - Giant Molecular Cloud
IMF - Initial Mass Function
SFR - Star Formation Rate
SFE - Star Formation Efficiency
KS law - Kennicutt-Schmidt law
DIG - Diffuse Ionised Gas
MW - Milky Way
PDF - Probability Distribution Function
SN - Supernova
AGN - Active Galactic Nuclei
AMR - Adaptive Mesh Refinement
SPH - Smoother Particle Hydrodynamics
RT - Radiative Transfer
RHD - Radiation Hydrodynamics
MHD - Magneto Hydrodynamics
CFL - Courant-Friedrichs-Lewy
NFW profile - Navarro-Frenk-White profile
A beginner’s guide to galaxies

Astronomers try to understand the Universe, and everything within it, through a combination of theoretical and observational approaches. This licentiate covers the subject of evolution and formation of disc galaxies, primarily, through the perspective of computer simulations. Current observations are not able to resolve the smallest scales of galaxies and, due to galaxies evolving on the scale of millions of year, can not directly study the evolution of galaxies. The computational, also called numerical, approach offers a more detailed view of the physics within galaxies and to more directly study the evolution of galaxies. A key concept discussed in this thesis is the connection between gas turbulence and other galaxy properties, such as star formation, over time. Our aim is to directly compare our results, from numerical simulations, with observational data that have investigated the same parameters.

From the most current and widely accepted cosmological framework theory, known as the ΛCDM model, astronomers believe that galaxies are formed in large and massive dark matter halos. Dark matter, while we do not know exactly what it is, interacts only via gravity and makes up 85% of all the mass in the Universe, which makes it crucial for galaxy formation. Each galaxy is encompassed by a dark matter halo and the properties of a galaxy is closely connected to the inherit properties of the halo it inhabits (such as mass and rotation, see Section 1.1). On a cosmological scale, dark matter is distributed throughout the Universe in an interconnected structure known as the "cosmic web", which helps funnel gas into these halos and aids galaxy formation. The first galaxies in the Universe formed roughly several hundred million years after the Big Bang and were formed by gas flowing into the gravitational potential wells produced by these massive dark matter halos. As gas accumulated, the cloud of gas started to compress into a disc-like structure and form stars; creating the first galaxies.

Galaxies are primarily made up from a combination of gas, dust, stars, and dark
matter. Disc galaxies, like our Milky Way, are the main type of star-forming galaxies, named from their distinct disc-shape of gas and stars. Commonly, disc galaxies have spiraling arms of gas and stars and a bulge of stars near the centre. However, there exists numerous types of galaxies with different morphologies; for example, some galaxies have a stellar bar extending from the central bulge, and the appearance of the spiral arms can vary drastically between disc galaxies. In this work, I will primarily use the umbrella term 'disc galaxies' to refer to all star-forming galaxies with discs, even though there is a sub-sample of diffuse or dwarf disc galaxies that form little to no stars. The gas in disc galaxies is referred to as the interstellar medium (ISM) and contains several distinct gas phases (see Chapter 2) with different temperatures and densities.

Astronomers know that stars are formed from huge gas clouds within galaxies, commonly confined in the spiral arms, and that galaxies then recycle their gas content to form new stars, but the intricate details of this process are not yet fully understood. Observations have shown that gas outflows are ubiquitous in star-forming galaxies and are part of the cycling of material known as the baryon cycle (see Section 2.2), which regulates the gas content of the galaxy. Many observational and numerical studies have shown that the gas within disc galaxies is highly turbulent and that turbulence imperative for several galactic properties, in particular star formation (see Section 2.4).

In this work, I have made numerical simulations (see Chapter 4 for details how simulations are made) in order to try to understand observations with highly turbulent interstellar medium. Numerical simulations have been successful in reproducing key properties of observed galaxies but several physical processes are usually omitted, due to how computationally expensive it would be to include all known physics. Understanding what physical processes drive observed correlations is crucial to learn from numerical simulations and then apply this knowledge to observations.

This thesis is outlined in the following way. In Chapter 1, I discuss the formation of the first galaxies in a cosmological context. In Chapter 2, I outline the gas phases within the interstellar medium and explain the formation of stars, various physical processes (particularly, star formation and stellar feedback), and go into depth about observations of the supersonic gas turbulence within gas discs. Chapter 3 regards the numerical side of galaxy research, particularly about the various methods employed in numerical codes. Finally, Chapter 4 contains a summary of the papers I have written and which this thesis is based on. I put my work into context of other similar work, as well as an outlook of
upcoming projects. This thesis contains two papers, which I will refer to as Paper I and Paper II throughout this thesis, and are available at the end of this thesis.
Chapter 1

Galaxy Formation and Evolution

High-redshift observational surveys have shown that the very first galaxies formed within the first few hundred million years of the Big Bang, corresponding to a redshift range of $z \sim 11 - 15$ (e.g. Coe et al. 2013; Oesch et al. 2016; Bunker et al. 2023; Austin et al. 2023). With the onset of the James Web Space Telescope (JWST), observations have been able to probe some of the first galaxies in the Universe. These observations have uncovered several massive galaxies, which are more numerous than expected from theory (given their high stellar mass) at this redshift ($z \gtrsim 15$; e.g. Boylan-Kolchin 2023). These discoveries highlight that there is yet more to be understood about the earliest galaxies and how they form and evolve. In the coming sections, I will discuss how galaxies evolve over time and how their properties are connected to the DM halo they form within.

1.1 Galaxies inherit properties from DM halos

The $\Lambda$CDM cosmology model is the most widely accepted framework for the evolution of the Universe, and predicts that galaxies form in massive Cold Dark Matter (CDM) halos which are self-bound with gravitational potential wells deep enough to accrete and condense gas. These dark matter (DM) halos were distributed in the Universe according to primordial oscillations that occurred in the very early Universe. This has been confirmed with the WMAP/Planck observations of the cosmic microwave background (see Springel et al. 2006, for a detailed review), from which we also learned that the amount of visible, baryonic, matter in the Universe only makes up $\Omega_b \approx 4.93\%$ of all matter in the Universe (Planck Collaboration et al. 2020). Additionally, the fraction of baryonic mass to total
mass is $\Omega_b/\Omega_m \approx 0.65\%$, which signifies that only a fraction of the total mass inside a galaxy is visible matter (gas, stars, etc). Furthermore, within galaxies, the DM halo seems to largely follow the same type of density profile (see Section 3.8).

No galaxy forms in complete isolation, and the environment they inhabit have a crucial impact on their evolution. The distribution of DM forms a cosmic web which facilitates gas to be funnelled into galaxies, but also creates nodes of galaxy clusters with varying sizes. In an isolated environment, disc galaxies can form when gas condenses into the halo (e.g. the first galaxies) and flattens into a disc due to the conservation of angular momentum. Accretion of gas into a disc galaxy is also important for the evolution of the disc, as it inserts angular momentum and more pristine gas to form stars.

As galaxies undergo interactions and various cluster-induced effects (gas stripping, galactic accretion, ram pressure stripping, ionising background, among others), their morphological structure tend to become less gas rich and spherical, like elliptical galaxies. This type of evolution can be inferred from the morphological distribution of galaxy populations in clusters, which have high density regions where galaxies often undergo interactions.
Thus, central galaxies are commonly massive elliptical galaxies while the galaxies on the outskirts are more likely to be disc galaxies. This is known as the morphology-density relation in galaxy clusters (first studied by Dressler 1980, but see e.g. Calvi et al. 2012 for a recent example).

Both observations of high-redshift galaxies and simulations agree that galaxies in the early Universe had irregular or clumpy structure, with intense star formation episodes. Thus, they were turbulent and had not yet formed a proper rotating disc, which is expected from ΛCDM theory. Observationally, an orderly rotating disc is commonly inferred from the ratio between rotational velocity and gas turbulence \( \frac{v_c}{\sigma_g} \). As galaxies undergo interactions or mergers, they become more turbulent and this ratio moves towards lower values.

Since disc galaxies evolve together with their halo, many of the properties of the gas and stellar disc of the galaxy can be related to the characteristics of the host halo. This technique of matching properties is commonly called abundance matching. For example, the rotational velocity of a galaxy can be related to the total mass of the halo through the virial theorem, which connects potential and kinetic energy in a self-gravitating system in equilibrium. Observationally, this is indirectly what yields the Tully-Fisher relation (Tully & Fisher 1977, but see also Straatman et al. 2017), which is an empirical law between galaxy luminosity (which is a proxy for stellar mass) and the rotational velocity in the outskirts of the gas disc.

Another example of abundance matching is the relation between halo mass and stellar mass, which allows astronomers to predict the typical amount of stellar mass given a halo mass. This relation is shown in Figure 1.1 for a specific redshift (local galaxies), since both the disc and halo mass evolve over time. Furthermore, theoretical work of disc galaxies have shown that many observational relations can be reproduced with analytical equations and a few assumptions of the halo and disc (e.g. Mo et al. 1998; Behroozi et al. 2013, 2019). Additionally, N-body simulations of DM halos have shown how e.g. DM density follow a specific profile. These kind of connections between the DM halo and the baryonic matter allows us to reduce the initial parameters when modelling entire disc galaxies. I will discuss some of these abundance matching techniques in more detail, and how they are implemented in my simulations, in Section 3.8.

In context of all the relations discussed above, our galaxy, the Milky Way (MW), is a very typical disc galaxy in several aspects. For example, most of the disc galaxies observed
in the local Universe have roughly the same halo mass as the MW (e.g. Behroozi et al. 2019). Furthermore, the MW lies along the main-sequence of galaxies (see Section 1.2) and contain many typical spiral galaxy features (e.g. spiral arms, stellar bar, supermassive black hole). Tracing a MW-like halo backwards in time, a MW-progenitor would be among one of the galaxies undergoing a starburst during the cosmic star formation peak, which is the topic of the next section.

1.2 Star formation in a cosmological context

Stars are an essential building block of galaxies and, as such, the formation of stars is an essential part of galaxy evolution. Stars inject heavy elements and energy into the interstellar medium through feedback (see Section 2.3) and regulate gas content (see Section 2.2). The first stars are believed to have formed in small DM halos, too small to sustain star formation and form into galaxies. As halos began hierarchically merging into larger and larger structures, gas rich halos began forming a disc and after some time the gas within the disc starts to clump up. High-redshift galaxies have been observed to have
Figure 1.3: The star formation rate versus the stellar mass of galaxies during the cosmic peak of star formation history, at redshift $\sim 2$. The data is coloured to separate between quiescent main galaxies and galaxies undergoing starbursts. Figure reproduced from Rodighiero et al. (2011).

a clumpy morphology and bursty star formation (e.g. Elmegreen et al. 2009; Zanella et al. 2019). However, the mass of these clumps might be overestimated in observations due to poor spatial resolution at high redshifts (e.g. Tamburello et al. 2017).

The Universe experienced a peak in cosmic star formation rate density, which is the amount of stars being formed within some large (cosmological scale) volume and time epoch, at a redshift of about $z \sim 2$. The evolution of this parameter can be seen in Figure 1.2. In the early Universe, galaxies contained mostly pristine, low-metallicity, gas which is very inefficient in cooling and, thus, forming cold dense clouds to allow for star formation. However, early galaxies contained massive gas reservoirs and were spatially closer and had more frequent interactions and mergers. This allowed individual galaxies in the very early Universe to form stars on the order of $\sim 100 M_\odot \text{yr}^{-1}$. Galaxies at redshift $z \sim 2$ still had an ample supply of cosmic gas inflowing, they were interacting, and the gas had been enriched with metals that increased the radiative cooling of the gas. The peak at $z \sim 2$ is
due to the combination of a ample supply of cosmic gas reservoirs, mergers between gas-rich galaxies being more common, and enough time had passed for halos to grow massive enough to produce high SFRs (see Madau & Dickinson 2014; Conselice 2014, for reviews on the evolution of SFRs and environment, respectively).

On the other hand, in the local Universe the most common galaxies, which have halo masses $10^{12} \, M_\odot$ (similar to the MW), have burnt through their gas supply and the accretion of new gas is minimal. In order for these galaxies to reach comparably high SFRs they need to be undergoing a merger, as this process delivers a fresh supply of gas and creates dense, gravitationally unstable, environments that boosts star formation through compressive turbulence (e.g. Renaud et al. 2009, 2014).

Galaxy formation is inefficient. As will be discussed further in Section 2.1, on a galactic scale disc galaxies only transform gas into stars at an efficiency on the order of $\sim 1\%$. Star formation, and thus galaxy formation, is thus a very inefficient process. Starburst galaxies have a much higher efficiency, but a starburst is a transient process which only covers a minimal time of a galaxy’s lifespan. This transient period can be understood through the galaxy main sequence (Daddi et al. 2007), which is defined through the tight relation between SFR-stellar mass as shown in Figure 1.3. Starburst galaxies lie at a significantly higher SFR for the same given stellar mass, i.e. above the main sequence. However, this above figure and explanation considers a redshift range $z \sim 1.5 – 2.5$ during which starbursts were a lot more common than today. Today, after this epoch of frequent gas accretion and merger, galaxies are more massive but also more quiescent, forming stars at a very low rate.
Chapter 2

The Interstellar Medium

The interstellar medium (ISM) is the mixture of gas and dust that fills the space between the stars in disc galaxies. The gas has many different phases with distinct temperatures and densities; for example, cold dense molecular gas, warm and cold neutral gas, and hot diffuse ionised gas. However, the ISM is a complex and widely dynamic system, with densities in the range $\rho_g \sim 10^{-4} - 10^4 \text{ cm}^{-3}$ and temperatures between $T_g \sim 10 \text{ K}$ and $T \sim 10^8 \text{ K}$ (see e.g. Ferrière 2001, for a review on the ISM). Majority of the ISM gas mass resides in cold, dense molecular clouds ($\rho_g \sim 10^2 - 10^6 \text{ cm}^{-3}$, $T \sim 10 - 20 \text{ K}$), which are confined to the spiral arms in disc galaxies or (in high-redshift galaxies) dense clumpy structures throughout the disc. Stars form from molecular clouds above some threshold density. Young stars are very luminous and their ionising radiation shape the environment around them, giving way to H II regions, which are warm, fully ionised regions inside molecular clouds ($\rho_g \sim 1 - 10^3 \text{ cm}^{-3}$, $T \sim 10^4 \text{ K}$). Filling most of the disc is neutral (mostly HI) gas of a wide range of temperatures ($\rho_g \sim 10^{-1} - 10^2 \text{ cm}^{-3}$, $T \sim 10^1 - 10^4 \text{ K}$). Warm, diffuse ionised gas (DIG; $\rho_g \sim 10^{-2} - 10^{-1} \text{ cm}^{-3}$, $T \sim 10^4 \text{ K}$), is spread throughout the disc as well as above and below (see Haffner et al. 2009, for a review on H II region and DIG). Surrounding the galactic disc, warm/hot ionised outflowing gas is suspended and, eventually, re-accreted into the galaxy. Details of how gas is recycled from stars back to the ISM is in Section 2.2.

Star formation is crucial for the formation of galaxies, but has been found to be very inefficient (see Section 2.1). Star formation requires the presence of molecular H$_2$ gas, due to its high density, cooling, and shielding from ionising UV radiation. Furthermore, dust grains, which are more complex molecular structures formed from metals, facilitate
the formation of $\text{H}_2$ by providing a nucleation site where atoms can cool. The ionised regions around stars, known as H II regions, are one of the main contributors of the H$\alpha$ emission line and is often used observationally to determine the star formation rate in galaxies. However, DIG, which fills much of the volume inside and above the galaxy disc, is also a significant contributor to the H$\alpha$ line. This line is prominent in observations as it comes from the recombination of the hydrogen atom between energy levels $3 \rightarrow 2$.

The heating mechanisms which ionise the gas come primarily from stellar feedback (see Section 2.3 for details) and are crucial in regulating the star formation in galaxies.

Finally, the ISM is supersonically turbulent in all of these phases; the cold molecular and the warm ionised gas phase alike. Analysing this turbulence, where it originates from, and comparing it with observations is the main focus of this thesis. Thus, I will give a detailed view of this subject in Section 2.4.

## 2.1 Star Formation

Stars form in groups, referred to as stellar populations, during the collapse of giant molecular clouds. The distribution of the initial mass distribution of the formed stars have been found to follow a similar function even at different star formation events. This initial mass function (IMF) has been observed through several methods (see e.g. Kroupa et al. 2013, for an overview): counting stars in the Milky Way, spectral energy distribution modelling, among others. The exact shape of the IMF, whether it evolves over time, or if it varies in extreme environments (such as near the galactic centre) is still under heavy discussion within the astronomy community (see e.g. Hopkins 2018, for a review).

Kennicutt (1989, 1998) analysed the gas surface density of neutral and molecular hydrogen gas, $\Sigma_g = \Sigma_{\text{HI}} + \Sigma_{\text{H}_2}$, in disc and starburst galaxies, and plotted it against the SFR density in each galaxy. This empirical correlation is what is now known as the Kennicutt-Schmidt (KS) law, after this work by Kennicutt and its initial discovery by Schmidt (1959). The law correlates the SFR density with the total gas density within a cloud as

$$\dot{\Sigma}_* \propto \Sigma_g^N.$$  \hspace{1cm} (2.1)

The value of $N$ depends on the environment, as local spirals with a quiescent star formation have $N \approx 1.0$ (Bigiel et al. 2008), but starburst have a larger index. For starburst galaxies,
Kennicutt (1998) calculated a value of \( N \approx 1.3 - 1.4 \), but starburst have been shown to have a different conversion factor between the observed CO molecule and molecular H\(_2\) gas, and correcting for this yields \( N \approx 1.8 \) (e.g. Narayanan et al. 2012). This law can be extended into a 3D law with a similar form, as I will discuss in Chapter 3 when implementing it in my simulations.

The seminal work of Bigiel et al. (2008) investigated this relation with sub-kpc observations of local disc and dwarf galaxies, and highlighted the efficiency of star formation in different types of galaxies. In Figure 2.1, the efficiency of local disc galaxies (the coloured contour) lies between 1 – 10%, while galaxies with more bursty star formation are in a transient period of transforming most of their gas into stars. However, at some cut-off density \( \Sigma_g \sim 10 M_\odot \text{ pc}^{-2} \) there is a large drop in efficiency. This threshold is believed to
signify a region in which the gas density is too low to boost the formation of molecular 
gas, which in turn forms stars. This is highlighted by the low-surface brightness (LSB, 
in the figure) galaxies that occupy this parameter space in the figure. Alternatively, the 
efficiency can be described in terms of depletion time $\tau_{\text{dep}} \equiv \Sigma_g/\Sigma_{\text{SFR}}$, which is the time 
for the galaxy to transform all of its gas into stars assuming the current SFR is constant. 
This yields an average $\tau_{\text{dep}} \sim 2$ Gyr for a non-starburst spiral galaxy with an efficiency of 
a few percent.

2.2 The Baryon Cycle

The baryon cycle in galaxies refers to the life cycle of the ISM, as gas is turned into 
stars and then re-processed into the ISM. This cycle involves the gas going through many 
different phase transitions, from molecular star-forming gas to ionised plasma, as well as 
gas fusing into heavier elements (inside stars) that then enrich the ISM with metals. The 
formation of stars is regulated by the stars themselves, through different stellar feedback 
effects (e.g. supernova, stellar winds, radiation; see Section 2.3) that heat and disrupt the 
formation of molecular clouds.

A cartoon of the baryon cycle can be seen in Figure 2.2. The figure shows a 2D 
histogram of the total gas density versus the temperature from one of my simulations. 
Annotated on top of the plot are arrows and numbers which, together with the figure 
description, describe the phases of the baryon cycle. What is omitted in this picture is 
the time-scale that gas spends in each phase. For example, gas is confined in molecular 
clouds for a long time before becoming stars. The time spent within stars varies widely 
and is dependent on the mass of the star; massive stars have volatile stellar winds and 
undergo SN swiftly, depositing their mass in the ISM, while low-mass stars live a long 
and quiet life. The ionised gas has many methods of cooling, which it does efficiently 
(see also Section 3.4 for how cooling is affected by temperature), but this depends on 
the environment; e.g. gas above/below the disc is too diffuse to cool effectively and is 
constantly heated by ionising radiation.
Figure 2.2: A temperature-density diagram of the gas from one of my simulations. Annotated on this diagram is the baryon cycle of material, illustrating how the gas is being continuously recycled. 1) Cold and dense molecular gas, part of giant molecular clouds, is the birth place of stars. As the large gas cloud is perturbed, it collapses and forms a multitude of stars. 2) Newly formed stars can warm and ionise its surroundings, causing the dense gas to be heated and ionised, forming H II regions. The gas in stars is used as fusion fuel in the core of the star and combines into heavier elements, but is also convected throughout the envelope. During a star’s lifetime, it ejects some of its gas envelope into the ISM, enriching it with warm and metal-rich gas. As a star undergoes supernova, most of its metal enriched gas is cast into the ISM (or out from the galaxy, forming outflows) at high speeds and temperatures. 3) Hot/Warm ionised gas cools down radiatively. 4) Warm gas keeps cooling through various processes until it forms dense molecular gas again.
2.3 Stellar Feedback

As mentioned in the previous section, stellar feedback is the process by which stars heat and enrich the ISM with metals, and is important for regulating the star formation within the ISM. The heating processes coming from supernovae (SNe), stellar winds, and ionising radiation are some of the primary mechanisms which form the variety in ISM phases. The momentum injected by supernova and, to a much lesser extent, stellar winds is believed to be partly responsible to the supersonic turbulence within the ISM (see Section 2.4).

Radiative feedback from young massive stars \( (M_\star \gtrsim 8 M_\odot) \) shapes the surrounding molecular clouds through very strong ionising radiation. This feedback effect is responsible for creating the, almost, fully ionised volume around the newly born stars within the molecular cloud, known as HII regions. These regions contribute a large fraction of Hα emission in a galaxy, as ionised hydrogen atoms recombine with electrons and cascades down the energy levels, until it (possibly) transitions from \( 3 \rightarrow 2 \) and emits an Hα photon with wavelength \( \lambda = 6563 \, \text{Å} \). Furthermore, radiation deposits a significant amount of momentum into the molecular cloud as first stars form, but stellar winds and supernova become the dominant feedback effects soon after.

Stellar winds occur due to complex dynamics caused by magnetic fields, radiation pressure, and convection inside the star, which ejects material from the gas envelope. In particular, massive stars can drive stellar winds powerful enough to shock into the ISM, which thermalises the wind and heats up the shocked material.

Furthermore, the most massive stars have short lifespans on the order of a few Myr, after which they undergo core-collapse supernova, also called Type II SN (SNII). The star then explodes, shooting very heavy material into the ISM and leaving only a neutron star or black hole, depending on the mass. Several, clustered, SNII can form expansive superbubbles of ionised gas that are much more explosive in nature. Stars which do not undergo SNII, but are massive enough to fuse carbon and oxygen end up ejecting most of their envelope, leaving only its core, and becoming white dwarfs. However, if the star has a companion from which it can accrete mass, it will eventually reach a threshold mass that causes the star to explode into a Type Ia SN (SNIa). This kind of supernova is more delayed than SNII, as it originates from a less massive star undergoing a slower fusion process before it is able to remove its envelope.

SNII and Ia are the most prominent types of SN and are therefore crucial astrophysical
mechanisms for shaping the ISM and driving hot galactic outflows. However, there are several more feedback effects that are important in heating the ISM and possibly drive turbulence. Black holes and neutron stars can drive feedback through (binary) accretion, creating a hot disc and magnetically driven jets of material. Supermassive black holes, which exist in the centre of most disc galaxies, can also accrete gas and drive very powerful feedback, which are then called active galactic nuclei (AGN). Cosmic rays, which are believed to originate from SN, AGN, and other effects, have also been found to have a big impact in shaping the ISM gas phases and structure; in particular, for heating the molecular gas and limiting star formation.

2.4 Gas Turbulence

Turbulence is the random motion within a medium, and a high turbulence level indicates that these motions have high speeds. This leads to regions of compression and high density when these velocities are converging, and diffuse gas when they are diverging. Thus, a highly turbulent medium can achieve higher densities than a medium with low turbulence. The ISM is supersonically turbulent, meaning that turbulence is a crucial component in shaping the properties of a galaxy, as will be further discussed in this section.

In an isothermal environment, where there is no heat exchange/injection within/into the gas, theoretical work has shown that the probability distribution function (PDF) of gas densities (i.e. the distribution of possible gas densities) is well-described by a log-normal distribution (see e.g. Kritsuk et al. 2007; Federrath et al. 2008, 2010; Salim et al. 2015). The standard deviation (width) of the density PDF scales with the Mach number $\mathcal{M}$, which is the factor of how much faster the gas is travelling than the speed of sound within that medium (i.e. $\mathcal{M} > 1$ is supersonic). Furthermore, the density PDF is dependent on other factors, such as the magnetic field strength, virial parameter $\alpha_{\text{vir}}$ (which indicates whether a system is gravitationally bound or collapsing), and the turbulence driving component (compressive versus solenoidal, see paragraph below). The log-normal density PDF is primarily relevant on a cloud-scale, and has been suggested to be the underlying parameter for mass distribution of molecular cores and stars (i.e. the core mass function and IMF, see e.g. Padoan et al. 1997; Padoan & Nordlund 2002; Hennebelle & Chabrier 2008), as well as the SFE and SFR (Federrath & Klessen 2012, 2013). However, the density PDF is still highly relevant at galactic scales, as it has been
applied to reproduce the KS-relation (Salim et al. 2015; Elmegreen 2018).

Furthermore, turbulence can be split into compressive and solenoidal components, which act radially (inwards/outwards) and rotationally, respectively. Solenoidal turbulence is a swirling turbulent flow that only stabilise the gas from collapse, while compressive gas can destabilise gas clumps and, thus, promote star formation. In a supersonically turbulent system, the components are, on average, in momentum equipartition (i.e. they contain the same amount of turbulent momentum). In numerical simulations, the shape of the density PDF has been found to vary strongly between these two turbulence components (e.g. Federrath et al. 2010). Additionally, simulations of galaxy merger have been shown to enchant the gas densities, and therefore star formation, due to an increase in compressive turbulence (Renaud et al. 2009, 2014).

Finally, it is not fully understood what the main driver of this turbulence is, but any process that injects energy and momentum into the ISM contributes to some degree. Observations show this correlation between turbulence and SFR on kpc-scales, thus processes connected to star formation are favoured; for example, two common sources discussed in literature are stellar feedback and gravitational instability. Power spectrum of the turbulent energy indicates that turbulence cascades, dissipationless, from larger scales to smaller until the motions dissipate into thermal (first predicted by Kolmogorov 1941, but see also Fensch et al. 2023 for a recent example), and so both local and global effects can drive turbulence at different scales inside a galaxy. Section 2.4.2 contains a more thorough discussion of the various theories investigated in literature.

2.4.1 Observations in the ISM

Supersonic turbulence is a ubiquitous feature of the ISM and is observed on a range of spatial scales, redshifts, and gas tracers (see e.g. Elmegreen & Scalo 2004; Mac Low & Klessen 2004; Glazebrook 2013, for reviews). Observationally, turbulence is commonly quantified as the full-width half-maximum (FWHM) of the observed line profile, while theoretical work often quantify it as the dispersion in velocities (see Section 3.9.2).

The level of turbulence observed in the neutral and molecular gas of local galaxies is of the order $\sigma_g \sim 10 \text{ km s}^{-1}$ while high-redshift galaxies are much more turbulent $\sigma_g \sim 50 \text{ km s}^{-1}$. However, the ionised gas can reach $\sigma_g \lesssim 40 \text{ km s}^{-1}$ in local galaxies and $\sigma_g \lesssim 100 \text{ km s}^{-1}$ in high-redshift galaxies. This $\sigma_g$ offset between gas phases is shown in
Figure 2.3 with a compilation of observed turbulence (y-axis) in local and high-redshift galaxies, split up into tracers of molecular, neutral, and ionised gas phases. This evolution of gas turbulence is expected from $\Lambda$CDM theory, as discs become rotationally dominated and less clumpy over time. However, there are also observational difficulties which can artificially enhance the observed turbulence, which I will detail in Section 2.4.3.

The turbulence within molecular clouds was first found to be closely tied to the size and mass of the clouds in the seminal work by Larson (1981). The mass-size and turbulence-size scaling relations in Larson follow $M \propto R^{2.0}$ and $\sigma \propto R^{0.5}$, respectively, and are similar to results from recent observations and simulations (e.g. Miville-Deschênes et al. 2017; Grisdale et al. 2018). This highlights that turbulence is a crucial component for molecular cloud properties, and thus star formation, at cloud-scale, $\sim 0.1 - 100$ pc. As discussed in Section 2.4, the gas density PDF is connected to supersonic turbulence, meaning that GMC turbulence (through the density PDF) affects the IMF and molecular cloud core properties, as well as the SFE and SFR (see Mac Low & Klessen 2004; Ballesteros-Paredes et al. 2007, for reviews on the relation between GMC turbulence and star formation).

However, this correlation between GMC turbulence and SFR also exists at galactic scales, as can be seen from observational data of the SFR-$\sigma_g$ relation in individual galaxies, presented in Figure 2.3. This connection between SFR and $\sigma_g$ is useful to compare underlying drivers of the ISM turbulence.

### 2.4.2 What Drives Turbulence in the ISM?

As mentioned above, gas turbulence is ubiquitous within many gas phases in the ISM, but the question remains where this turbulent energy is coming from. Turbulence from internal, thermal motions is expected to contribute at most a few km s$^{-1}$ in GMCs (this can be calculated using Eq.3.17, assuming an ideal gas). Many different origins have been suggested and tested in literature, e.g. stellar feedback, gravitational instability, rotational shear, gas accretion (primarily at high-redshift), magnetic field interaction, AGN (see Mac Low & Klessen 2004; Elmegreen & Scalo 2004, for reviews). While all of these astrophysical processes likely inject some amount of turbulence, not all of them could sustain it over a longer period of time. Currently, there is no consensus among astronomers which effects contribute the most, but two origins have been heavily discussed in literature: stellar feedback och gravitational instability.
Figure 2.3: A compilation of observations, ranging from local to high-redshift galaxies, of the turbulence (quantified as the line-width, $\sigma_g$) and star formation rate (SFR) in the ISM of galaxies. Most of the data points represent individual galaxies and the data is divided into the different gas phases the observations derived $\sigma_g$ for. Left: The $\sigma_g$-SFR relationship as depicted in most literature. Right: Same as left, but with a logarithmic y-axis to highlight how the scaling (red lines) of the relation changes around $\text{SFR} \sim 1 \text{M}_\odot \text{yr}^{-1}$. Figure is from Ejdetjärn et al. (2022); see their Table A1 for data references.

Stellar feedback provides a natural bridge between turbulence and SFR, as recent star formation is closely connected to strong outbursts of kinetic energy (in the form of supernova or stellar winds). The energy ejected by SNe has been argued to be sufficient to drive turbulence in gas-rich disc galaxies, where SNe are frequent enough to sustain an input of turbulent energy larger than the dissipation (Hayward & Hopkins 2017).

Gravitational instability is also very relevant, due to its close connection to density, which both star formation (KS law, Section 2.1) and turbulence (density PDF, Section 2.4) are closely tied to. The turbulence in disc galaxies has long been thought to be correlated with gravitational instability. Toomre (1964) derived an analytical equation for a parameter $Q$ that governs the stability of a self-gravitational disc in hydrostatic equilibrium. The expression for $Q$ connects mass density $\Sigma_g$ with the velocity dispersion $\sigma_g$, at some
radius $R$ in the disc, as

$$Q_g = \frac{\sigma_g \kappa}{\pi G \Sigma_g}$$

(2.2)

where $G$ is the gravitational constant and $\kappa$ the epicyclic frequency. The stellar $Q$ parameter has approximately the same equation. When $Q < 1$ for the entire disc, it becomes unstable and self-gravity causes it to collapse in on itself. However, as a gas disc collapses, the density increases and the gas exerts more outward pressure/turbulence, stabilising the disc. Although it is also possible that stellar feedback is boosted due to an SFR boost as gas density increases. Therefore, $Q = 1$ is considered a quasi-stable state in which the disc is self-regulating by collapsing then increasing gas turbulence and re-stabilising, which might also be a natural evolution of discs galaxies. This model has been further developed in e.g. Romeo & Falstad (2013), who developed a method for combining $Q$ of several components, in particular stars and gas of different phases. Comparing different $Q$ components shows how stars are driving gravitational instability in local discs with a low gas fraction.

In Krumholz et al. (2018), the authors present an analytical model (based on their previous model Krumholz & Burkert 2010), derived from considering a quasi-stable state gravitational disc in hydrostatic equilibrium. The model features heavy parameterisation and has been widely used to interpret galaxy observations of $\sigma_g$, which can be used for deriving properties of the observed environment. The author argues that gravity is the dominant driver of turbulence, but with stellar feedback being more relevant in some cases. However, these models are focused on the bulk motion of the gas and do not represent the turbulent motions of diffuse (ionised) gas, which complicates comparisons with observational data using Hα emission as a tracer for disc dynamics.

### 2.4.3 Observational Artifacts

Extra-galactic observations of the ISM need to take into account several effects in order to create a consistent picture of how certain parameters, such as the velocity dispersion, evolve with redshift. Even if only considering disc galaxies at the same redshift, the range of masses, sizes, morphology, etc. that these galaxies could have is quite large. Furthermore, the environment within and surrounding these galaxies varies sharply between the local and high-redshift galaxies. Thus, these factors have to be corrected for in order to attain comparable data between observations and simulations.
Relevant to observations of turbulence, beam smearing is a mechanism that artificially broaden emission lines due to a combination of insufficient spatial resolution and steep velocity gradients. The poor resolution causes the steep velocity gradient to be unresolved and the velocity field to be "smeared", which leads to broadening of emission lines. In observations of the ISM, the rotational velocity curve is a prominent contributor to beam smearing and must be subtracted from the velocity field. This involves analytical or numerical modelling of the disc, which in turn requires a constraint on the inclination by which the galaxy is being observed. In literature, several methods have been developed to correct for this, for example direct 2D morphological modelling or 3D fitting from datacubes (e.g. GALFIT, 3D-Barolo Peng et al. 2002; Di Teodoro & Fraternali 2015), but no method is perfect and the modelling is further complicated by lack of data points to fit (due to poor spatial resolution) for high-redshift galaxies. Furthermore, inner disc regions are more affected by beam smearing due to the sharp velocity gradients of the rotational curve near the centre.

Observations commonly try to represent resolved galaxies with one, global, $\sigma_g$ value, which is accomplished by performing a weighted average of $\sigma_g$ in each element/pixel. The choice of weights can have a contribution on the global $\sigma_g$, as flux-weighting has been suggested to favour the inner regions of discs (e.g. Davies et al. 2011; Di Teodoro & Fraternali 2015), which are more turbulent but also more prone to be affected by beam smearing (e.g. Józsa et al. 2007).

Additionally, the turbulence in high-redshift galaxies has mostly been observed in Hα (see Figure 2.3), due to being a ubiquitous line and its strength. As Hα traces ionised gas, $\sigma_g$ should be compared with other observations of ionised gas, since molecular and neutral gas phases are much less turbulent (see Section 2.4).
Chapter 3

Numerical Modelling of Galaxies

Hydrodynamical simulations are a popular tool in astrophysics for studying the gas in a variety of environments; for example, galaxy formation, stellar formation, supernova explosions, planetary discs, planet atmospheres, among others. There are several different numerical techniques used to perform these simulations, each with its own strengths and weaknesses (see Springel 2010b; Teyssier 2015, for reviews). Essentially, hydrodynamic codes apply fluid equations (see Section 3.2) to simulate the motion of the gas and an $N$-body approach to calculate the movement of particles (e.g. stellar, DM).

There are many complex physical processes occurring within disc galaxies (see e.g. the feedback discussion in Section 3.6), and the impact of each process depends on the scale at which they are simulated. Physics that is not resolved is implemented through sub-grid recipes. Thus, numerical modelling of galaxies can be considered to be done on three distinct scales, which aim to investigate different phenomena and each category has its own advantages and drawbacks. Cosmological simulations model large-scale environment starting from assumptions of the initial matter density fluctuations, which then evolve to form more and more massive DM halos which, eventually, form galaxies. This approach is useful for describing the evolution of galaxies with the cosmic web and external influences such as galaxy interaction, mergers, and gas inflow/outflow. However, galaxies in cosmological simulations are not well-resolved and in order to study a particular galaxy in detail requires to single it out among the other galaxies in the large simulation volume and then re-simulate the sub-volume at a higher resolution. This is known as a zoom-in simulation and is a useful follow-up that allows the study of chosen galaxies in a realistic cosmological environment.
The other type of simulation are isolated galaxy simulations. As these do not include the effects of the surrounding environment, the drawback is that they only track the secular evolution of the galaxy, which is realistic for local field galaxies or for evolving galaxies over a short time-span during which external effects may be less relevant. However, a big advantage of this method compared to cosmological simulations is the reduced need of computational power, which facilitates the simulation to be run at much higher resolution for the same computational resources. Furthermore, an isolated galaxy can be modelled more explicitly according to the conditions that are desired for the research topic.

The final approach is to study individual patches within galaxies. This allows for a more detailed view of the complex physics occurring inside the ISM at small scales and does not rely on as many sub-grid recipes as simulations of entire galaxies does. Additionally, the results assists in forming sub-grid recipes for galaxies at larger scales.

### 3.1 Numerical Codes and Physics

In this work, I use the hydro+$N$-body code RAMSES (Teyssier 2002), which is a popular code within the astrophysics galaxy community and has been part of large-scale projects such as: the cosmological SPHINX simulations (Rosdahl et al. 2018), the VINTERGATAN collaboration (focusing on a MW-like galaxy from zoom-in simulations; Agertz et al. 2021), and EDGE (considering dwarf galaxies; Agertz et al. 2020). Gas is treated as a fluid, following the Euler equations (Section 3.2), and stars and DM are treated as collisionless particles. RAMSES has several modules that allows for simulating physical effects such as magnetic fields and radiative transfer, but only the latter is of interest in this thesis.

Explicit implementation of physical processes yield a more realistic treatment, as it avoids sub-grid recipes, but it also requires more computational resources or to reduce the simulation resolution. For this reason, simulations commonly consider only a set of physics (including sub-grid) that are relevant for the research topic and researchers consider e.g. radiation and magnetic fields as secondary effects. Simulations with radiation hydrodynamics (RHD) apply radiative transfer equations on-the-fly and can therefore follow the propagation and evolution of spectra, but are also better at tracing the ionisation structure in galaxies. Magneto hydrodynamics (MHD) trace the magnetic fields of gas and objects, which then affects the motion of charged particles (within the gas) through e.g. braking (which can slow down rotational and turbulent motions). Magnetic fields
have been found to act as a stabilising force for molecular clouds, but are also important in the ionised gas to drive e.g. Alfvén waves, outflows, synchrotron radiation, and cosmic rays. However, high-resolution MHD simulations of isolated galaxies by Su et al. (2017) indicated that on, galactic scale, magnetic fields have a negligible effect on regulating the SFR, outflows, and balancing gas phases (see also Krumholz & Federrath 2019, for a review on magnetic fields’ impact on the SFR and IMF). This does not dismiss the importance of magnetic fields in galaxies, but indicates that it is a secondary effect for reproducing global trends in galaxies.

Many different numerical codes have been developed that are often used within the community, e.g. AREPO (responsible for the Illustris and IllustrisTNG cosmological simulations; Genel et al. 2014; Vogelsberger et al. 2014; Marinacci et al. 2018), GIZMO (used within the FIRE projects; Hopkins et al. 2014), GADGET (Millenium simulations, GRIFFIN; Springel et al. 2005; Lahén et al. 2020), among many others. Throughout the coming sections, I will go over numerical methods and physics implementations relevant for the RAMSES code and briefly put them into context of other numerical codes.

### 3.2 Hydrodynamic Equations

The numerical simulation of gas in galactic environments follow the same hydrodynamic equations that describe the conservation of mass, momentum, and energy in fluids. These equations can take on slightly different forms depending on the field in which they will be applied, but are independent on the method of implementation. In this work, we consider the Euler equations, which are a subset of the Navier-Stokes equations that do not consider the fluid viscosity. This approximation works well in most astrophysical applications, as gas is not very viscous, but depending on the field of research and application (see the discussion on SPH in Section 3.3) it may be non-negligible and an "artificial viscosity" term needs to be added. The Euler equations are as follows,

\[
\frac{\delta \rho}{\delta t} + \nabla \cdot (\rho u) = 0, \tag{3.1}
\]

\[
\frac{\delta \rho u}{\delta t} + \nabla \cdot (\rho u \otimes u) + \nabla p = -\rho \nabla \phi, \tag{3.2}
\]

\[
\frac{\delta \rho e}{\delta t} + \nabla \cdot [\rho u (e + p/\rho)] = -\rho u \cdot \nabla \phi, \tag{3.3}
\]
where $\rho$ is the total density, $u$ is the velocity vector, $e$ is the internal (thermal) energy of the gas, $p$ is the pressure, and $\phi$ is the source term for the gravitational potential of the gas and other objects (e.g. stars). The first equation captures the conservation of mass (and is known as the continuity equation), the second is momentum, and the third is the energy.

The internal properties of the gas need to be connected to each other through some equation of state. Commonly, in galactic-scale simulations, the ideal gas law is applied:

$$p = n k_B T = \frac{\rho k_B T}{\mu m_p} = (\gamma - 1) \rho e$$

(3.4)

where $\gamma = C_p/C_V$ is the adiabatic index, defined as the ratio between specific heat capacity at constant pressure $C_p$ and constant volume $C_V$. For a mono-atomic gas, e.g. neutral hydrogen, the adiabatic index is $\gamma = 5/3$ and for a diatomic gas, e.g. molecular hydrogen, it is $\gamma = 7/5$. Within the ISM, most of the volume is mono-atomic and, thus, simulations commonly apply $\gamma = 5/3$.

The gravitational potential, both internal and external, can be solved through the Poisson equation

$$\nabla^2 \phi = 4\pi G \rho_{\text{mass}},$$

(3.5)

where $\rho_{\text{mass}}$ is the density of all the matter (typically that involves gas, stars, plus DM). Calculating the total gravitational acceleration induced on one gas cell/particle individually between every other cell/particle is computationally expensive. Instead, an efficient technique is usually employed to only calculate the individual acceleration between objects that are close-by (within the same volume), while acceleration of objects far away are calculated using the total combined mass density within a larger volume. In RAMSES, the total mass density within grid cells for particles is calculated using a cloud-in-cell interpolation (see e.g. Hockney & Eastwood 1981), which treats particles as a cubical mass distribution and inject its mass onto surrounding cells according to the volume overlap between the 'cloud' and simulation cells.

As will be discussed in Section 3.6, when implementing stellar feedback thermal energy, momentum, and mass are injected into the surrounding gas and are treated self-sufficiently by these equations. However, some additional terms need to be added when introducing more physics, e.g. magnetic fields and radiative transfer. In particular, RHD simulations need to add a non-thermal energy term to represent the ionisation and heating associated with the radiation field.
Numerical codes track the motion of the gas by solving for the acceleration from internal and external forces, e.g. pressure and gravity. Depending on the numerical method employed, e.g. SPH or AMR (see the section below), the framework of solving these equations differ significantly. In the case of a mesh-based approach, the code must solve the Riemann discontinuity (cell boundary) problem to calculate the flux between mesh cells. One method of doing this is following the Godunov scheme, which is a finite-volume framework for numerically solving partial differential equations with boundary conditions (i.e. Riemann problems). For example, RAMSES employs a 2nd order Godunov scheme (see Teyssier 2002, for a detailed explanation), which reconstructs the distribution of hydro variables (e.g. density, pressure, velocity) in each simulation grid cell as a piece-wise slope. Then, the scheme calculates the flux (through the hydro equations) at each boundary of adjacent cells. Finally, the quantity in each cell is updated by the flux out/in of the cell by integrating over the chosen time-step.

3.3 Adaptive Mesh Refinement

Numerical codes apply the hydrodynamic equations on the gas elements within a simulation volume. How to trace the gas elements is a choice that differs between numerical codes and can be divided into either following the gas as it moves through space (Lagrangian; also called mesh-free) or observing it at fixed points of a mesh (Eulerian). The results from codes with different methods have overall been in agreement (see Section 3.6), but each method has inaccuracies/uncertainties that need to be addressed within the code. Here I will go over the basic idea behind each method but with focus on AMR, as it is used in RAMSES.

One of the prominent mesh-free methods is Smoothed Particle Hydrodynamics (SPH; Lucy 1977; Gingold & Monaghan 1977; Monaghan 1992, see also Springel 2010b for a recent review) framework, available in e.g. GADGET (Springel et al. 2005, 2021), GASOLINE (Wadsley et al. 2004), AREPO (Springel 2010a), GIZMO (Hopkins 2015). This approach represents each gas element as a particle with the relevant hydro parameters, i.e. mass, velocity, and internal energy. The particles have a smooth kernel function which represents the size/volume distribution of the gas element and can vary between particles. An issue with traditional SPH methods is their inability to fully capture shocks and fluid-mixing instabilities (Agertz et al. 2007). To address this, artificial viscosity and
other recipes are added in modern SPH codes.

In a mesh-based approach each gas element is represented with the hydro parameters within a mesh, commonly with a quadratic grid structure. However, having a grid of constant size throughout the simulation volume would be computationally expensive for high-resolution simulations and so Adaptive Mesh Refinement (AMR) techniques are instead employed. The AMR method works like a grid but allows high-resolution in individual grid cells that are of interest, according to some user-defined criteria. Despite this, the contact discontinuities between grid cells are believed to cause angular momentum loss which could be significant over a longer period of time. A visualisation example of AMR, from one of my RAMSES simulations, is presented in Figure 3.1. Example of some astrophysical codes with AMR framework are FLASH (Fryxell et al. 2000), ENZO (Bryan et al. 2014), RAMSES.

Furthermore, codes with mixed implementations that combine Lagrangian and Eulerian approach are becoming more popular and are continuously being developed and improved upon. Unstructured moving mesh methods form a mesh, e.g. through Voronoi tessellation, that deforms at each time step to contain a finite mass or volume, and are employed in e.g. AREPO, GIZMO, TESS (Duffell & MacFadyen 2011). A possible issue with these codes are numerical diffusion of the gas between cells. Another type of mixed implementation is the Lagrangian meshless method in GIZMO, which traces the gas as particles but calculate the hydro equations on the boundaries (constructed mesh) created from kernel "overlap". These mixed codes aim to employ the strengths of each approach, but they are relatively new and it remains to be understood if they inherit any of the drawbacks as well (see the review on cosmological simulations by Somerville & Davé 2015).

### 3.3.1 Refinement Strategies

The RAMSES code utilises an AMR technique, which works by refining only the grid cells that have been marked for refinement according to some user-defined criteria. Refinement is then performed by splitting up the parent cell into more (child) cells with less mass and fewer particles. Each sibling cell together are called an oct. A refined cell is commonly chosen to be half the size (along each axis) of its parent cell, which means that, for a three-dimensional volume, eight child cells are made inside the parent cell. The hydro
Figure 3.1: A gas density slice map with the AMR structure overplotted. The data is from a RAMSES simulation output and shows a slice of the ISM inside a gaseous disc. The regions with the highest gas density have the most resolved AMR cells.
and $N$-body motions are solved simultaneously, for each refinement level, starting with the most refined gas cells.

The time-steps for integrating the gas motion is commonly synchronised by choosing the time-step of a child cell to be half as large as its parent cell. This means that the parent cell is synchronised with its child cell every other time-step, which avoids erroneous time-interpolation and the associated instabilities. The time-steps are chosen as the smallest time-step required by various local conditions of the gas cells. For example, the Courant-Friedrichs-Lewy (CFL) condition is a maximum time-step constraint applied in most numerical codes as a base condition. The criteria concern the propagation of density waves and requires the time-step to be small enough for the wave not to cross (jump over) an entire simulation cell, which would cause various instabilities and artifacts in the numerical solution. The CFL conditions is also used for other quantities, e.g. for radiation in RHD. In the case of RT, the time-step is commonly a time sub-step chosen to be synchronised with the hydro time-step. Furthermore, in RAMSES, to ensure that the transition in spatial scale between cells is smooth, the difference in refinement level between adjacent cells can not exceed one refinement level. Following this methodology, when a cell is refined its neighbouring cells are also marked for refinement.

A common strategy is to refine based on the mass inside a cell, in order to capture the fragmentation of dense gas regions. Additionally, by refining the mesh with the strategy to keep the amount of particles inside a cell constant, the mesh structure follows the dynamics of the particles (a sort of quasi-Lagrangian approach). Beyond the criteria described here, there are many more that could be defined according to what environments/physics the simulator wants to capture.

A noteworthy aspect of astrophysical simulations is that they are computationally expensive and typically require supercomputers with hundreds of cores over the duration of several months. Therefore, numerical codes have to be written and optimised to distribute and synchronise computations from several cores at once. RAMSES uses a Hilbert space mapping to transform the 3D cell mapping to a 1D list, which can then efficiently distribute the list of cells between the available cores. To keep track of the refinement level of cells of the AMR structure, RAMSES applies a double-linked hierarchical data structure known as a "Fully-Threaded Tree" (Khokhlov 1998), which keeps track of which child cell belongs to which parent cell, and vice-versa. When a cell is refined, another entry is added to this list, and when it is de-refined the entry is removed.
3.4 Gas Cooling and Heating

Gas cooling and heating are crucial to represent the gas phase transitions inside the ISM (see Section 2.2 for details). Cooling is primarily done by emitting radiation and allows gas to cool down into cold molecular gas and then form stars, while heating can be done by e.g. stellar feedback and creates the warm and hot gas phases of the ISM. The amount of radiative cooling is dependent on the interactions between charged particles within the medium, which is set by the ionisation conditions and thermal motions of the medium. In turn, the amount of ions and internal motions is related to the gas temperature and amount of metals (and dust) available. A high metallicity facilitates gas cooling, since metals have more electromagnetic transitions and can therefore absorb ionising photons and emit a cascade of photons with less energy (but the same total energy) that can escape the ISM; this is known as emission line cooling. In simulations without RT, ionisation populations are calculated through the Saha equation, which assumes a collisional ionisation equilibrium and only depends on density and temperature (i.e. no ionising radiation from stars).

In simulations, the cooling is commonly represented with a cooling function, which dictates the rate of heat loss as a function of temperature and metallicity. When implementing a cooling function, it is generally assumed the gas is optically thin, i.e. most of the radiation will escape. An example of cooling functions used in my simulations is presented in Figure 3.2 (check the figure description for technical details), for different gas metallicities. As can be seen, the gas has different regimes of cooling, with a stark transition at $T \approx 10^4$ K, which is the characteristic temperature associated with hydrogen recombination lines. The temperatures below this are mainly from fine structure cooling, while the cooling above comes from metal lines emission (e.g. Helium recombination). Brehmsstrahlung radiation, from freely interacting ionised particles, is the main responsible effect for the tail at $T \gtrsim 10^7$ K.

There are numerous astrophysical effects that heat the gas, which are implemented with the various sub-grid recipes (see e.g. the stellar feedback approach in Section 3.6). In the case of RHD, stars explicitly emit radiation that heats the gas (see Section 3.7 for how cooling/heating with RT is implemented). Furthermore, a UV background is often adopted to represent the UV heating from the surrounding environment present since the epoch of reionisation (see e.g. the approach by Faucher-Giguère et al. 2009).
Figure 3.2: Example of cooling functions used in my RAMSES simulations. The function appearance at $T > 10^4$ K comes from metal cooling tables from Sutherland & Dopita (1993). Gas with $T \leq 10^4$ K use the fine structure cooling rates from Rosen & Bregman (1995). Figure is reproduced from Agertz et al. (2013).
3.5 Star Formation Recipe

As discussed in Section 1.2, stellar properties are connected to the parent cloud from which they are born. Observationally, star formation follows the KS law and, assuming a standard IMF, thousands of stars will form with a wide range of masses. Capturing the formation of all stars as individual particles in galaxy-scale simulations is usually not feasible due to the extensive computational time and memory required; although, some recent simulations have been able to trace the most massive stars as individual particles (see e.g. the GRIFFIN project Lahén et al. 2020). Instead, during a star formation event a star particle is formed that represent an entire stellar population following a standard IMF.

A generalised, three-dimensional version of the KS law (see Eq. 2.1) can be written as

\[ \rho_* = \epsilon_{\text{ff}} \frac{\rho_g}{t_{\text{ff}}} \], for \( \rho > \rho_* \),

where \( t_{\text{ff}} = \sqrt{3\pi / 32 G \rho} \) is the free-fall time for a spherical molecular cloud and \( \epsilon_{\text{ff}} \) is the star formation efficiency for the collapse of a spherical cloud. The density threshold \( \rho_* \) for star formation to occur can be chosen arbitrarily, as long as it reflects the typical density of molecular clouds where star formation occurs. A common choice is \( \rho_* \approx 100 \text{ cm}^{-3} \), as most of the gas above this density is molecular (see Gnedin et al. 2009, and references therein). Numerical work has been able to reproduce the low efficiency \( \epsilon_{\text{ff}} \sim 1\% \) observed in local disc galaxies (when averaged over kpc scales; see Figure 2.1) by setting \( \epsilon_{\text{ff}} \approx 10\% \) (see e.g. Agertz & Kravtsov 2015, 2016; Grisdale et al. 2018, 2019).

At each star formation event, the above star formation equation is stochastically sampled following a discrete Poisson distribution, which decides how many stellar particles would form during this event. The probability of \( N \) events occurring is

\[ P(N) = \frac{\lambda_P}{N!} \exp(-\lambda_P) \]

where \( \lambda_P \) is the average rate of the events. Then, one stellar particle is formed with mass equal to the total mass of the stars that would be formed. Each stellar particle corresponds to a stellar population with mass distribution following a standard IMF. The choice of IMF mainly impacts the amount and nature of stellar feedback in the simulations, as it governs the distribution of massive stars. In my simulations, a Kroupa (2001) IMF is
applied with the form

$$\Phi(M) = A \times \begin{cases} 2M^{-1.3}, & \text{for } 0.1 \leq M < 0.5 \, M_\odot \\ M^{-2.3}, & \text{for } 0.5 \leq M < 100 \, M_\odot \end{cases}$$

(3.8)

where $A$ is chosen to normalise the total mass to be equal to the stellar particle mass $M_\star$. Compared to more bottom (low-mass) heavy IMFs (e.g. Kroupa et al. 1993), the Kroupa (2001) IMF produces over twice as many core-collapse SN, while a Chabrier (2003) IMF produces roughly three times as many. Keep in mind, the chosen IMF is normalised to fit the feedback budget and so most IMF can be used. Note that the IMF is here extrapolated to include stellar masses upwards of $100 \, M_\odot$. Furthermore, the canonical Kroupa IMF extends to lower masses, but the feedback contribution from this mass range is negligible.

3.6 Stellar Feedback

Simulations of the ISM have in the past decade focused on implementing only a few stellar feedback processes, due to the limits of computational resources and time constraints. Owing to recent advancements of numerical codes and computational resources, large-scale projects with detailed physics models have started to emerge. However, even the most highly resolved ISM-patch simulations need to employ some type of sub-grid recipes. Some of the feedback processes included in galactic-scale simulations include: individual and clustered SNe, stellar winds, photoionisation, radiation pressure, cosmic rays, stellar binary evolution, black hole feedback (see e.g. Somerville & Davé 2015, for a review of numerical simulations).

Particularly, the effects of these processes that needs to be modelled in simulations is their imparted momenta, mass, metallicity, and thermal energy. Additionally, RHD tracks the non-thermal energy and ion species, while MHD needs to account for the magnetic field created by these processes. Relevant to the work presented in this thesis, I only consider internal gas physics, gravity, stellar feedback, and explicit RT calculations.

The occurrence of SNe events can be tracked through stellar evolution models. The rate of SNe can then be expressed as a function of stellar age and metallicity (see e.g. Raiteri et al. 1996; Mannucci et al. 2006). At each timestep, the function stochastically probes whether a stellar particle contain any stars undergoing (individual) SN. Kim &
Ostriker (2015) showed that in order for a SN to be resolved in a numerical simulation, the cooling radius of the SN needs to be resolved by at least six cells. A SN that is resolved will simply input its energy ($10^{51}$ ergs) as thermal into its host cell, and the shockwave is treated self-sufficiently by the hydrodynamics. An unresolved SN assumes the shock has reached the Sedov-Taylor phase (the initial SN ejecta has swept up ISM material and no longer dominates the shock), and the resulting thermal energy and momentum is imparted into the host and surrounding cells. The metal enrichment from a typical SN can be calculated through hydrodynamics and explosive nucleosynthesis models (e.g. Thielemann et al. 1986, for SN Ia and Woosley & Weaver 1995, for SN II).

Similarly, stellar winds eject mass, energy, and metals into the ISM and the amount can be derived through stellar evolution models. However, stellar wind shocks shells are commonly on much smaller scales and, thus, are usually described with sub-grid recipes. Strong winds, from massive stars ($M \gtrsim 5 M_\odot$), will shock into the surrounding ISM and thermalise, transforming a substantial amount of its initial kinetic energy into thermal. Furthermore, despite not contributing significantly to the feedback energy budget, less massive stars ($M \gtrsim 8 M_\odot$) eject a significant amount of mass and metals into the ISM which needs to be modelled.

Radiation dominates the momentum injection early on after a stellar birth. In order for radiation to impart momentum into the gas, it has to be absorbed/scattered. Gas with higher metal contents (or, equivalently, dusty regions) have more electromagnetic transitions, which makes it more opaque to a wide range of frequencies. Thus, metal-rich gas is more likely to experience strong radiation pressure. In the case of RT not being calculated explicitly, the injection of momenta from radiation pressure has to be calculated analytically with a sub-grid recipe as

$$\dot{p}_{\text{rad}} = (\eta_1 + \eta_2 \tau_{\text{IR}}) \frac{L(t)}{c},$$

(3.9)

where $\tau_{\text{IR}}$ is the optical depth of infrared radiation, $\eta_1$ represents the radiation from the star while $\eta_2$ represents the contribution of re-processed IR radiation. $L(t)/c$ is the total momentum of the photons from the stars, where the $L(t)$ is the bolometric luminosity of the stellar population and $c$ is the speed of light.

The implementation of stellar feedback that we follow is outlined in Agertz et al. (2013). This approach traces the thermal energy, momentum, and metallicity ejection coming from stellar winds and supernovae (type II and Ia). Radiation pressure is also accounted for,
but for simulations where radiative transfer is explicitly calculated, momentum and energy is self-sufficiently tracked by the RT module (see section below). The stellar feedback effects considered are the (thermal) energy, momentum, and metal enrichment from SNe type Ia & II, stellar winds, and radiation pressure. This routine tracks the metallicity of iron and oxygen as separate hydro variables, due to these metals making up 90% of the total SNII ejecta mass (Agertz et al. 2013). Furthermore, oxygen in the ISM originates mainly from type II SNe and is an $\alpha$-element, which effectively makes it a proxy for other $\alpha$-elements. RAMSES uses the combined, total, metallicity for calculating the cooling function (see Section 3.4). The metallicity is transported with the bulk motion of the gas and is calculated as

$$Z = 2.09 \times Z_O + 1.06 \times Z_{Fe},$$

which assumes the gas follows a relative solar abundance metallicity for the $\alpha$-elements and iron (Asplund et al. 2009).

However, the specifics of stellar feedback implementation varies between numerical works and rely on empirical relation from observations as well as theoretical stellar synthesis models. The RAMSES code has been applied as part of the AGORA collaboration (Kim et al. 2014, 2016; Roca-Fàbrega et al. 2021), which had the goal of unifying and comparing various simulation codes (RAMSES, GADGET, ENZO, GIZMO, AREPO, GEAR, CHANGA, ART-I) by their results of MW-like galaxy simulations. The first and third paper of the project focused on cosmological zoom-in simulations while the second paper focused on an isolated MW galaxy. The project results showed an overall agreement between codes when applying a common physics model between the simulations, which validates the approach and results of papers applying similar physics for different codes. Similarly, code comparisons made by Hu et al. (2023) of an isolated dwarf galaxy also showed an overall agreement between RAMSES, GIZMO, AREPO, GADGET. However, cosmological simulations in the Aquila project (Scannapieco et al. 2012) showed significant differences in the galactic properties between codes, but attributed it to the variations in feedback implementation between the codes. Regarding cosmological simulations of galaxy formation, Somerville & Davé (2015) suggested codes are largely in agreement. And Naab & Ostriker (2017) noted in their review of galaxy simulations that discrepancies between codes and physics implementations have now been consolidated, by saying:

"The fact that many models — even if conceptually very different — driving a reason-
able galactic wind can successfully reproduce galaxy abundances and disk galaxy morpholo-
gies [...] indicates that the essential characteristics of the problem have been disclosed."

3.7 Radiative Transfer

Radiation hydrodynamic (RHD) simulations of galaxies have become more frequent in
the last decade, as astronomers realise the importance of RT in research topics focused on
the ionisation condition of the gas, e.g. the re-ionisation epoch or HII regions. However,
propagation of radiation on-the-fly is computationally expensive due to the number of
dimensions required; position, angular direction, and emission frequency. Several types
of numerical methods have been developed to efficiently couple the RT equations with
hydrodynamics, but can in essence be split up into two methods. Ray-based schemes
cast bolometric rays from each source, and update the radiation intensity and local gas
properties as it scatters throughout the simulation volume. There are several methods
of simplification, e.g. merging rays or approximating rays as cones. Moment-based RT
reduces the angular dimensions by deriving a simplified RT equation that considers only
the angular moments. In this section, I will describe the moment-based method applied
in RAMSES-RT, but the specifics of the RT suite is explained in Paper II.

In particular, RAMSES-RT uses the M1 (first moment) closure approach, which con-
nects the third moment of radiation (pressure) with the first one (energy) through an
Eddington tensor (see e.g. Levermore 1984). A moment-based method is diffusive in na-
ture, meaning that the radiation can be thought of as a bulk of field or fluid "diffusing"
throughout the gas. The simplifications of the RT equation then results in conservation
laws, akin to the Euler equations. This allows this method to be readily coupled to the
AMR grid and the flux be integrated with a method similar to the Godunov scheme (see
Section 3.3 for an explanation on how this work for hydrodynamics). Specifically, the
radiation flux between the cells is calculated with the Global Lax-Friedrichs method (see
Rosdahl et al. 2013), which is similar to the Godunov scheme but avoids the Riemann
problem; the boundary interaction can be reasonably ignored with the M1 closure ap-
proach, as the photon bulk flow does not self-interact in the same way as a gas does.
Furthermore, to make the time-steps (and thus computational time) manageable, the
speed of light is commonly reduced by two orders of magnitude, which is viable due to
the reduced speed still being much faster than the propagation speed of the gas and/or
Each star emits radiation in wave-packets with a range of frequencies, each packet having a characteristic (average) energy and intensity. By clever choices of wave-packet frequencies, the impact of radiation stellar feedback can be represented with only a few wave-packets. Some of the physical processes that are vital to capture with RT are: disassociation of H$_2$ molecules from UV photons (also called Lyman-Werner photons), radiation pressure from single scattering, heating through photoionisation, and non-thermal IR radiation from multi-scattering. The intensity of each wave-packet is calculated from spectral energy distributions of each particle, assuming a single stellar population.

Since radiation heats the gas, a non-thermal energy source term needs to be added to the hydrodynamic equations (see Eqs. 3.1) and hydro variables. The cooling, formation, and destruction of molecular hydrogen H$_2$ is important for star formation and needs to be described with an explicit recipe (see e.g. Nickerson et al. 2019) while H$_2$ self-shielding can be handled self-consistently within the RHD. Furthermore, the ion species that are to be traced need to be implemented as additional hydro variables, and should be related to what the wave-packet frequencies are tracing (e.g. Lyman-Werner photons requires H$_2$ as a species).

### 3.8 Initial Conditions

The simulations presented in this work are all isolated disc galaxies, constructed to be Milky Way-like galaxies, in terms of their mass and size, at different redshifts. I model these galaxies using various density profiles for the initial gas, stellar, and dark matter distributions. The exact values of the initial conditions used in my simulations are all presented in Paper I and II, but here I will provide a broader view that would be excessive in a research paper.

The spin, concentration, and shape of dark matter halos have been constrained by N-body simulations of DM halos, as they evolve over time in a cosmological ΛCDM environment. Navarro et al. (1996) first showed that DM halos dominantly follow, what is now known as, a Navarro-Frenk-White (NFW) profile of the shape

$$\rho_h(r) = \frac{\rho_0}{r/r_h(1 + r/r_h)^2}, \quad (3.11)$$

where $\rho_0$ is a constant that varies between halos. The mass of a halo can be assumed
to be close to the virial mass, which is the mass required for an over-density to be self-
gravitationally bound and in virial equilibrium. Furthermore, it is often assumed to be
200 times the mass at the critical density $\rho_c$ of the Universe, which is the cosmological
density required for a flat Universe and the virial mass is then defined as

$$M_{\text{vir}} = M_{200} = 200 \frac{4\pi \rho_c r_{200}^3}{3}. \quad (3.12)$$

In these equations $r_h = r_{200}/c$ is the halo scale length, $c$ is the dark matter halo concen-
tration and $r_{200}$ is the virial radius. Other profiles, such as the Einasto profile (Einasto
1965), have been suggested to better match certain DM halos due to the fact that it, like
some observed DM halos (primarily those of dwarf galaxies), have a constant density at
the core while the NFW is divergent.

Through the same numerical method, Macciò et al. (2007) showed that galaxies at
redshift zero have a DM concentration related to their mass as

$$\log c = 1.020 \pm 0.015 - 0.109 \pm 0.005 (\log M_{200} - 12), \quad (3.13)$$

where $h \equiv H_0/(100 \text{ km s}^{-1} \text{ Mpc}^{-1})$ is the reduced Hubble constant and $H_0$ is the Hubble
constant.

The stellar and gas disc are treated as following an exponential surface density like

$$\Sigma(r) = \Sigma_0 \times \exp(-r/r_d), \quad (3.14)$$

where $r_d$ is the scale height or radius and $\Sigma_0$ is the central density which normalises
the expression such that integrating yields the total mass. For a typical disc galaxy, the scale
radii and heights of the gas and stellar component are usually similar but with the gas
disc being more extended in height and radius.

The stellar bulge follows a spherical Hernquist potential of the form

$$\rho(r) = M_b \frac{r_b}{2\pi r (r + r_b)^3}, \quad (3.15)$$

where $r_b$ is the scale radius and $M_b$ the total mass of the bulge.

In order to, in practice, determine the initial position and velocity of gas and particles
in our simulations, I use the code GALIC (Yurin & Springel 2014). The code constructs
$N$-body models following density distributions and initial guesses for the velocity, which it
iterates over to find values in which they system is in a collisionless dynamical equilibrium.
For the DM halo, GALIC takes the initial virial mass, rotation, and concentration as input parameters. Using the parameterisation of DM halos by Mo et al. (1998), I calculate from the desired virial mass the expected halo properties. The initial stellar and gas properties were based on those used in the AGORA project for an isolated MW galaxy (Kim et al. 2016), but varying the gas fraction to represent the MW at different redshifts. The gas is initialised as particles but these are only used to form the AMR mesh, and then the mass of each gas particles is deposited in their corresponding grid cell.

3.9 Calculating Observational Parameters

Due to the different nature between simulation volumes and observational data, it is often not possible to calculate parameters in the exact same way between the two. For example, stellar populations in simulations have a definitive mass, but observational data have waveband or spectral data that is used to infer the stellar mass. RAMSES, like all other numerical codes, creates data outputs at times specified by the user. The data outputs contain the hydro and N-body parameters of each simulation cell, from which all analysis is done. Here, I will give a general explanation of how to calculate these parameters but the specifics are available in Paper I and II.

3.9.1 Star Formation Rate

Simulation outputs of stellar particles contain, for each particle, information about its age and mass, among other parameters. Calculating the SFR history can then simply be done by binning the stars by age and calculating the difference of stellar mass between each bin. The choice of bin size is important for comparisons with observational SFRs. Observationally, the SFR is calculated by assuming that the SFR is constant during some period of time and the exact time depends on the SFR tracer used. For example, UV comes from young massive stars and is thus instantaneous, with a lifetime of a few Myr, while Hα in H II regions fades away after \( \sim 5 - 10 \) Myr (see e.g. Krumholz 2015, for an overview).
3.9.2 Velocity Dispersion (Turbulence)

Observational methods commonly quantify turbulence as the line width (FWHM) of a profile line (see Section 2.4). In theoretical work, turbulence is viewed as the dispersion of velocity of the objects (e.g. stars or gas) analysed, calculated as the weighted standard deviation,

$$\sigma = \sqrt{\frac{\sum w_i (v_i - \bar{v})^2}{\sum w_i}},$$  \hfill (3.16)

where $v_i$ is the velocity of one element, $w_i$ is the weight of one element, and $\bar{v}$ is the mean velocity. Weighting is important to represent $\sigma_g$ in the bulk of gas by using density/mass as weights, but is also important to trace specific gas phases/tracers. Additionally, velocity dispersion can also be defined as the square-root of the second moment (variance) of the velocity.

When calculating $\sigma_g$ in a simulation volume, it does not take into account the internal motions of the gas in each simulation cell. The internal speed of the gas can be derived for an ideal gas (i.e. following Eq.3.4) as

$$c_s = \sqrt{\frac{k_B T}{\mu m_p}} = \sqrt{\frac{P}{\rho}}. \hfill (3.17)$$

The total turbulent motions is then achieved by adding these together in quadrature,

$$\sigma_g = \sqrt{\sigma_{g,0}^2 + c_s^2}.$$

As mentioned in Section 2.4, resolved observations commonly calculate the $\sigma_g$ in each observed region to a global weighted-average $\sigma_g$. Similarly, simulations can calculate $\sigma_g$ locally and then combine to a global value. Unresolved observations observe the total line profile of the galaxy, which is essentially the same as calculating the velocity dispersion of all the gas in inside and around the galaxy.

Simulations are able to produce mock profiles by binning the desired quantity, e.g. mass or luminosity, in velocity bins and sum the quantity in each bin. From this, the line-width of the distribution can be determined, which yields a closer equivalent to the observational method. However, this adds further complications as some line profiles might be complex and not have a straight-forward determination of the line-width.
3.9.3 Hα Luminosity

The emissivity of a specific emission line inside a gas can be described with the equation

$$\epsilon_{\text{H} \alpha} = n_e n_H h \nu q,$$

where the electron number density $n_e$ and hydrogen number density $n_H$ of the gas, $\nu$ is the light frequency, $h$ is the Planck constant and $q$ is the emission rate. The work presented in these papers concerns the ionised gas tracer Hα, which is the $3p \rightarrow 2s$ (Balmer-α) transition in hydrogen with transitional energy $h\nu = 3.026 \times 10^{-12}$ ergs. There are more hydrogen fine-structure lines from the third excitation level down to the second (e.g. $3d \rightarrow 2p$) but the transition rate of these are comparably negligible.

Ionisation of the HI gas can occur either collisionally or radiatively and the total emission rate is a combination of these. The collisional component is calculated as

$$q_{\text{coll}}(T) = \frac{1.3 \times 10^{-6}}{T^{0.5}} \left( \frac{T}{11.2} \right)^{0.305} \times \exp \left( \frac{-\Delta E}{kT} \right),$$

which is a fit of observational atomic data from Callaway et al. (1987). The recombination rate is given by Dijkstra (2017) as

$$q_{\text{recom}} = \epsilon_{\text{H} \alpha}^B(T) \alpha_B(T),$$

where $\epsilon_{\text{H} \alpha}^B(T)$ is the Hα emission rate and $\alpha_B(T)$ is the probability that the recombination of electron and proton, called case B recombination, will cascade through the $3p \rightarrow 2s$ transition and form an Hα photon. These are calculated as

$$\epsilon_{\text{H} \alpha}^B(T) = 8.176 \times 10^{-8} - 7.46 \times 10^{-3} \log_{10} \left( \frac{T}{10^4} \right) + 0.45101 \left( \frac{T}{10^4} \right)^{-0.1013},$$

where the shape is fitted using observational values from Storey & Hummer (1995), and

$$\alpha_B(T) = 2.753 \times 10^{-14} \left( \frac{315614}{T} \right)^{1.5} \left( 1 + \left( \frac{315614}{T} \right)^{0.407} \right)^{-2.42}$$

is a fit from Hui & Gnedin (1997).

Furthermore, as I perform mock observation of Hα, it is relevant to consider HII regions, as they correspond to a large fraction of the total Hα emission. The size of HII regions are commonly quantified as Strömgren spheres with radius

$$R_{SG} = \left( \frac{3 N_i}{4 \pi \alpha n_H^2} \right)^{1/3},$$

38
where $N_i$ is the rate of ionising photons, $\alpha$ is the radiative recombination coefficient. This radius represents the fully ionised region ($n_e = n_H$) from calculations of the ionisation balance between UV radiation from the central star and recombination of the gas.

As an aside, in Paper I, Eq. 3.19 has $\Delta E$ replaced with the H\textalpha{} photon energy $h\nu$. This stems from an inaccurate assumption that the hydrogen atom is in a constant state of excitation to the 2p energy level, which is incorrect in most environments as most of the H\textsc{i} gas is in its ground state. The impact of this could be noticeable in environments dominated by collisional de-excitation.
Chapter 4

Summary and Context of My Work

The work presented in this thesis focuses primarily on the gas turbulence within the ISM of disc galaxies, using numerical simulations. There have been numerous hydrodynamical simulations in the past that have investigated gas turbulence in the ISM of disc galaxies (e.g. Hung et al. 2019; Orr et al. 2020; Kohandel et al. 2020; Jiménez et al. 2022; Kohandel et al. 2023; Rathjen et al. 2023). These are a mix of simulations looking at cosmological, isolated galaxy, and ISM patches. My work adds to this by investigating the relationship between gas turbulence $\sigma_g$ in different gas phases, with a focus on direct comparisons with observational data. Paper I focused on the observed relation between $\sigma_g$ and SFR (seen in Figure 2.3) and exploring the various observational artifacts underlying this data. Paper II focused on the commonly used Hα ionised gas tracer in order to determine which environments within, and outside, the galaxy that make up the observed line profile. Both papers present simulations of isolated Milky Way-like galaxies at redshifts corresponding to $z \sim 0$ and $z \sim 2$, but Paper II employs RHD. A more thorough summary and context of these paper follows below.

4.1 Paper I

In this work, I investigated the correlation between gas turbulence $\sigma_g$ and star formation rate (SFR) inside my simulated disc galaxies. By performing mock observations to mimic observational data, I was able to perform a more direct comparison with my simulations and evaluate the impact of observational effects, such as beam smearing (from resolution and galaxy inclination) and the choice of gas tracers. Furthermore, I performed the simu-
lations with and without stellar feedback effects, to understand its role on ISM turbulence. Observations have previously shown that SFR-$\sigma_g$ follows a piecewise power law (see Figure 2.3), but this data is very heterogeneous and contains galaxies with widely different properties, at different redshifts, observed at different resolutions and inclinations.

I found that inclinations, if not (accurately) corrected for with disc modelling, have a significant impact on the observed $\sigma_g$. This issue is primarily important for high-redshift galaxies with a high SFR and (presumably) high $\sigma_g$, since they are barely resolved by observations and thus the modelling becomes inaccurate. The resolution of the mock observations was found to have a minor impact on $\sigma_g$, unless observing galaxies near edge-on. Finally, the choice of gas tracer was found to be very important, as ionised gas (traced by my H$\alpha$ mock observations) was on the order of 2-3 times more turbulent than the neutral and molecular gas.

Additionally, removing stellar feedback from the simulations proved to not make any significant difference on the bulk motions of the gas, i.e. $\sigma_g$ of all the gas. However, without stellar feedback the ionised gas phase within the simulations was much weaker, and so was $\sigma_g$ of the ionised gas. Thus, feedback is important for driving turbulence in the ionised gas.

4.2 Paper II

In Paper II, I explored the origin of ionised gas and performed mock observation of a specific ionised gas tracer, H$\alpha$, to conduct more direct comparisons with observations. The initial conditions and setup of these simulations are almost identical to those in Paper I, but for this project we decided to include on-the-fly radiative transfer equations. This is important to capture the HII regions within the galaxy, which are one of the dominant sources of H$\alpha$ emission.

My results showed that despite the simulated galaxies having comparable outflow, SFR, and $L_{\text{H$\alpha$}}$ to observed galaxies, the contribution of H$\alpha$ emission from just outside the gas disc was negligible. Thus, the H$\alpha$ profile line represents the dynamics from inside the gas disc, but this does not take into account outflows that had yet to leave the disc. Furthermore, the fraction of H$\alpha$ from diffuse ionised gas (DIG) versus HII regions was on the low end of what has been observed, meaning that H$\alpha$ was dominated by HII regions. However, similar simulation by Tacchella et al. (2022) found a much larger contribution,
which is likely due to their post-processing of the emission line, which allows for the H\(\alpha\) line to scatter away from H\(\text{II}\) regions.

Furthermore, I showed that H\(\text{II}\) regions and DIG are distinctly different in the H\(\alpha\) luminosity and \(\sigma_g\) space. The \(\sigma_g\) of H\(\text{II}\) regions is around 2-3 times lower than for the DIG, which might reflect the environment (dense molecular gas) which H\(\text{II}\) regions inhabit. This is related to one of the next projects concerning the \(\sigma_g\) discrepancy between molecular and ionised gas.

## 4.3 Upcoming projects

I have two primary project remaining that will be part of my PhD dissertation. The simulations for these projects have already been done and a preliminary analysis has been made, but more analysis is needed and there are some aspects of these simulations that might need adjustment.

### 4.3.1 Haro 11

The next project I will finish up is to numerically model the Haro 11 dwarf galaxy, which is the closest confirmed Lyman-continuum galaxy and is believed to currently be undergoing a major merger. Furthermore, it has been thoroughly studied in most wavebands and is a respected member of the galaxy group at Stockholm University. Simulations can aid in interpreting all the observational data. Recent observations of the H\(I\) 21-cm line with the MEERKAT instrument revealed a long, \(\sim 40\) kpc, tidal tail of neutral gas coming from the galaxy (Le Reste et al. 2023). This, along with it undergoing a starburst, stellar population properties, and complex kinematics, suggest that it is undergoing a merger which is shaping its current properties.

The goal is to understand a possible path of origin for Haro 11 and to explain some of its current properties. Related to the other projects, I will also analyse the turbulence of this system in order to understand how a merging, starburst galaxy compares to the isolated galaxies presented in this thesis. Numerical modelling specific galaxies are uncommon, especially for dwarf-galaxies, and this kind of numerical work has never been done for Haro 11 before. This project might be split up into two papers, one focused on comparing the global properties of my simulated galaxy to Haro 11, while a follow-up paper would
include RT to analyse how radiation is escaping the knots of star formation within the galaxy.

Related to the Haro 11 project, I will model the ESO-338 galaxy, which also has a tidal tail and been suggested to be undergoing a merger (Cannon et al. 2004). This project involves investigating new HI data of this tidal structure, which is currently being gathered with MEERKAT, but my primary contribution will be to perform the simulation of ESO-338, which will aid with interpretations of the observational data.

### 4.3.2 Gas turbulence in molecular vs ionised gas

The final project I have planned is also about the ISM turbulence in isolated disc galaxies, but focusing on an offset in $\sigma_g$ between molecular and ionised gas that was highlighted by Girard et al. (2021). They showed that the ionised gas is roughly 2-3 more times as turbulent as the molecular gas and that this offset was constant for a wide range of galaxy gas fractions. Indeed, this was something that I observed in Paper I but has also been reported in simulations of individual ISM patches (Rathjen et al. 2023). In Paper II, I showed that the diffuse ionised gas has 2-3 times larger $\sigma_g$ than the dense ionised gas; similar to what has been observed between the ionised and molecular gas. This hints towards the diffuse gas driving this offset. For this upcoming project, I will investigate this offset and whether it is universal or varies in different local environments within the galaxy.
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