Formation and Evolution of Early-Type Galaxies. III Dependence of the star formation history on the total mass and initial over-density

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ABSTRACT

We investigate the influence of the initial over-densities and masses of proto-galaxies on their subsequent evolution (the star formation history in particular), to understand whether these key parameters are sufficient to account for the varied properties of the galactic populations. By means of fully hydrodynamical N-body simulations performed with the code \textsc{Evol}, we produce twelve self-similar models of early-type galaxies of different initial masses and over-densities, and follow their evolution from the early epochs (detachment from the linear regime and Hubble flow at $z \geq 20$) down to the stage when mass assembly is complete i.e. $z \leq 1$ (in some cases the models are calculated up to $z = 0$. The simulations include radiative cooling, star formation, stellar energy feedback, re-ionizing photo-heating background, and chemical enrichment of the interstellar medium; we do not consider the possible presence of Active Nuclei. We find a strong correlation between the initial properties of the proto-haloes and their subsequent star formation histories. Massive ($M_{\text{tot}} \approx 10^{13} M_\odot$) haloes experience a single, intense burst of star formation (with rates $\geq 10^3 M_\odot$/yr) at early epochs, consistently with observations, with less pronounced dependence on the initial over-density; intermediate mass ($M_{\text{tot}} \approx 10^{11} M_\odot$) haloes have histories that strongly depend on their initial over-density, whereas low mass haloes ($M_{\text{tot}} \approx 10^9 M_\odot$) always have erratic, bursting like star forming histories, due to the “galactic breathing” phenomenon. The model galaxies have morphological, structural, and chemical properties resembling those of real galaxies, even though some disagreement still occurs, likely a consequence of some numerical choices. We conclude that total mass and initial over-density drive the star formation histories of early type galaxies. The model galaxies belong to the so-called quasi-monolithic (or early hierarchical) scenario in the sense the aggregation of lumps of dark and baryonic matter is completed very early on in their history. In this picture, nature seems to play the dominant role, whereas nurture has a secondary importance.

Key words: Galaxies, cosmology

1 INTRODUCTION

Understanding how early-type galaxies (ETGs) formed and evolved along the Hubble time is still a formidable challenge. According to the current view of the subject, galaxies probably began to form at $z \sim 20 - 50$, when Dark Matter (DM) haloes containing Baryonic Matter in cosmological proportions gave origin to the first sufficiently deep gravitational potential wells (Tegmark et al. 1996). However, it is not yet clear how the paradigm $\Lambda$CDM cosmological model (which prescribes a hierarchical, bottom-top formation of structures, implying that massive systems assemble their mass later than the smaller ones) could be reconciled with the observational evidence for large, and red, galaxies already in place at very high redshifts (see e.g. Marchesini et al. 2011, Harrison et al. 2011, Mortlock et al. 2011). The problem is twofold: on one hand, it must be explained how and when star formation is quenched in both massive and small haloes, which is necessary to reconcile the theoreti-
cal prediction with the observed galaxy mass function (see e.g. Bundy et al. 2006); on the other hand, it should be also clarified how such massive systems can form at very high redshifts. Semi-analytical models of hierarchical formation (e.g., De Lucia et al. 2004, De Lucia & Blaizot 2007, Almeida et al. 2007, Gonzalez et al. 2009, Parry et al. 2004, De Lucia et al. 2011) have succeeded in explaining some of the observational features of ETGs, but fail in others. In particular they overestimate the number of faint objects and underpredict the sizes of bright ETGs. In addition to this, the hierarchical formation finds it difficult to explain the trend shown by the so-called α-enhancement at varying the mass of the galaxy. While massive ETGs are enhanced in α-element, the low mass ones are not. This requires that the history of star formation varies with the galaxy mass in a way that can be hardly met by the hierarchical scheme. However, since the topic has been examined in detail by (Tantalo & Chiosi 2002) and no new arguments have been made available we will not discuss it in details (but see Sect. 3.3).

All the above issues are essentially a reformulation of the long-standing dichotomy between the so-called hierarchi- cal and monolithic models of galaxy formation, as described respectively e.g. by Eggen et al. (1962) and by White & Rees (1978), possibly including some more sophisticated features (e.g. the dry merger scenario, as proposed by Bell et al. 2004). After the Cold Dark Matter cosmological model has got worldwide acceptance, the hierarchical model appeared to be the most convincing scenario of galaxy formation. In recent times, however, the situation has become by far more intriguing than in the past. Evidences claiming for a new scenario very close indeed to the classically and theoretically (see e.g. Chiosi & Carraro 2002), but recently they have become compelling (e.g. Thomas 2011; Gobat et al. 2011).

Merlin & Chiosi (2006, 2007) suggested that a series of primordial mergers of small stellar subunits (in a sort of early hierarchical galaxy formation mechanism), taking place at z ≥ 2, could lead to the formation of massive objects in the early Universe, resembling the observed systems as far as many different aspects (morphology, density profiles, metallicity...) are concerned. This picture finds support in numerical simulations of the very first stellar populations in the early universe, such as those by Bromm et al. (2002) and Yoshida et al. (2003), who found that the first virialized objects might form at z ∼ 30 with a typical mass of 10^6 M_☉. In this scheme, a first generation of stars in clumps inside primordial haloes of DM would first enrich the medium in metals, and also act as the building blocks of larger systems with masses up to 10^12 M_☉ via gravitational clustering. The mode of star formation and the efficiency of it is likely driven by the mass and mean density of the proto-galaxy as long ago pointed out by Chiosi & Carraro (2002) and more recently argued on observational basis by Harrison et al. (2011).

A fundamental piece of the puzzle is the role played by energy feedback during the formation of the galactic systems. Although it has long been known (see e.g. McKee & Ostriker 1977) that supernova explosions can inject large amounts of kinetic and thermal energy into the galactic Interstellar Medium (ISM), only more recently the importance of this phenomenon as a source of pressure sup-
the possibilities of the monolithic (or early hierarchical) mode of galaxy formation we have just outlined and simultaneously to bridge a gap between the results presented by Chiosi & Carraro (2002) and those by Merlin & Chiosi (2006, 2007). In the first study, the initial conditions for the proto-galaxy were not derived from self consistent cosmological simulations, but were created ad-hoc to model a spherical lump of DM + BM (in the right proportions depending on the adopted cosmological background) subject to collapse: the DM particles were distributed following the Navarro et al. (1996) universal profile, with velocities derived from the velocity dispersion \(\sigma(r)\) for a spherical, isotropic, and collision-less system with the adopted density profile (Binney & Tremaine 1987); while the BM particles (in form of gas at the beginning) were distributed homogeneously inside the DM halo with zero velocity field, mimicking the infall of primordial gas into the potential well of DM (White & Rees 1978). From this stage, star formation in the proto-galaxies was allowed to occur. Finally, supposing that at each redshift the over-density of proto-galaxies could vary within a suitable interval, many models were calculated at varying the total mass and initial over-density.

In the two studies by Merlin & Chiosi (2006, 2007), the initial conditions for DM and BM particles of the proto-galaxies were derived from large scale cosmological simulations, however adapted to our purposes. Since the procedure is much similar to the one adopted here, they will be jointly described below. However, in those papers no attempt was made to confirm the Chiosi & Carraro (2002) scenario about the mass and/or initial density dependence of the star formation history of a galaxy.

With this target in mind, the NB-TSPH code EvoL was fully revised, re-written in parallel language and much improving the input physics. All details about the physical ingredients and the many tests we have performed to validate the code can be found in Merlin (2009); Merlin et al. (2010) to whom the reader should refer. The present paper completes the analysis initiated with Chiosi & Carraro (2002) and continued with Merlin & Chiosi (2006, 2007). Here, we focus on the relation between the initial conditions of the host halo and the final properties of the galaxy making use of the parallel code EvoL and a consistent cosmological scenario to set the initial conditions. We investigate the process leading to the formation of isolated galaxies with different total masses and initial over-densities and cast light on whether the internal properties of a “real”, cosmologically consistent galactic halo made of DM and BM are sufficient to obtain the variety of structural and physical properties which characterize the population of early-type galaxies and to confirm the predictions made by Chiosi & Carraro (2002).

Furthermore, we will examine whether the size predicted by the new models better agree with the observational data to confirm the predictions made by Chiosi & Carraro (2002). Moreover, we will examine whether the internal properties of a “real”, cosmologically consistent halo should be generally considered part of its nature, since it is essentially fixed by the cosmological perturbations from which the halo forms, rather than part of the nurture processes, as it is sometimes viewed.

By varying the initial halo over-density of our models (in the way we are going to describe below), we try to mimic the action of the “natural” over-density. In some way, one could consider our “dense” models as systems belonging to generally over-dense regions ultimately becoming groups or clusters of galaxies, and our “less dense” models as totally isolated (field) galaxies.

All the model galaxies are obtained from the same initial proto-halo, however suitably modified in order to systematically change its properties from model to model (see below). We also run a few more cases with different prescriptions, to better explore the effects of some key physical quantities and assumptions.

The paper is subdivided as follows. In Sect. 2 we describe the numerical methods and the setting of the initial conditions. In Sect. 3 we analyze the results of the simulations. In Sect. 4 we discuss the mass density profiles of the stars we have obtained for the models. In Sect. 5 we present a preliminary discussion of the Mass-Radius relationship for ETGs. Finally, in Sect. 6 we draw some final remarks and conclusions.

**2 INPUT PHYSICS OF THE NB-TSPH MODEL GALAXIES**

**2.1 Setting of the initial conditions**

The ideal procedure to derive the initial structure of a proto-galaxy would be to start from large scale cosmological NB simulations of typically ~500 Mpc on a side in the framework of a given cosmological model so that the appearance, growth and subsequent aggregation of perturbations of all...
scales can be suitably described, cfr. e.g. the Millennium Simulation described in [Springel et al. 2005] who studied the formation, evolution and clustering of DM haloes in the Λ-CDM cosmology. Of course these simulations require a huge number of particles; for instance, the Millennium run followed more than 10 billions particles. Considering the huge number of DM haloes (proto-galaxies candidates) that come into existence, only a small number of particles will be used to describe the internal structure of many of them, the smaller ones in particular. If this way of proceeding is fully satisfactory from the point of view of a large scale cosmological simulation, it is not viable in the case we want to study in detail a single object of galactic size. This indeed will have total masses ranging from $10^8$ to $10^{13} \ M_\odot$, dimensions from say 1 to several tens of kpc, $M_{DM}/M_{BM}$ in cosmological ratio where DM is by far dominating, and to be properly described will require a large number of particles (at least 50,000) for each component. Repeated zooming in of smaller sub-portions of the whole initial grid is often applied when interested to small size objects such as individual galaxies compared to the cosmological volume. However, this procedure is quite expensive in terms of computational resources. Other strategies can be found that have already been adopted in some of the studies we referred to. In particular, an isolated perturbation can be artificially created and evolved without following the evolution of the much larger region of space in which it is located. Moreover, since the aim of this study is to explore a wide range of masses and initial densities, the creation of an ad-hoc set of haloes is the most appropriate method to adopt, instead of searching in large scale simulations for the particular haloes suited to our purposes. To this aim, we proceed as follows.

i) We assume the Λ-CDM concordance cosmology, with values inferred from the WMAP-5 data [Hinshaw et al. 2009]: flat geometry, $H_0 = 70.1 \ km/s/Mpc$, $\Omega_\Lambda = 0.721$, $\Omega_b = 0.046$ (giving a baryon ratio of $\approx 0.1656$), $\sigma_8 = 0.817$, and $n = 0.96$.

ii) To describe the growth of primordial perturbations, we use COSMICS, the free software written by [Bertschinger 1995]. This is particularly easy to use and it has already been adopted in many studies as the generator of initial conditions for NB simulations. However, we are not interested in a full cosmological simulation containing perturbations at all scales, but only in a portion of it containing a perturbation with properties (over-density, mass, and dimensions) fixed a priori. The size of this sub-section is fixed in such a way that the wavelength of the perturbation corresponding to the chosen over-density and mass is similar to (but suitably smaller than) the size of the sub-section. Our aim is to construct a reference proto-halo containing DM and BM particles in the right proportions each of which with its own known mass, position and velocity vector. This reference proto-halo, instead of being constructed by hand as in [Chiosi & Carraro 2002], is obtained by means of COSMICS in the context of an assigned cosmological scenario. Casting the problem in a different way, instead of searching within a large scale realistic cosmological box the perturbation most suited to our purposes, we suppose that a perturbation with the desired properties is already there, and derive the positions and velocities of all its DM and BM particles from a self-consistent, small-size cosmological box tailored to the perturbation we have chosen.

iv) This reference halo will correspond to the largest mass of the whole sample of galaxy models we have calculated. The grid containing our density perturbation has the size of $l = 9.2$ comoving Mpc on a side, and it is populated by a grid of $46^3$ particles. The regular positions of the particles are perturbed by COSMICS consistently with cosmological random gaussian fluctuations. Furthermore, a density peak is constrained to form at some very early epoch a virialized structure near to the center of the box (COSMICS allows us to specify the properties of the constrained density peak to model the density field in the desired way). We impose a gaussian spherical over-density with average linear density contrast $\delta \rho = 3$, smoothed over a region of radius $3.5$ comoving Mpc. We remind the reader that the linear density contrast is defined as $< \rho > /\rho_{bg} - 1$ (where $\rho_{bg}$ is the average matter density of the Universe), and it can be extrapolated beyond the linear regime; it would be equal to 1.86 at the epoch of virialization. Requiring a higher value in COSMICS simply bounces back in time the epoch of virialization.

v) Starting from these input assumptions, COSMICS returns the initial comoving positions and the initial peculiar velocities of the particles at the time in which the highest density perturbation is exiting the linear regime (i.e. it has $\rho/\rho_{bg} - 1 = 1$). This is an important point to remember. COSMICS is used only up to the exit from the linear regime. The subsequent evolution of the DM and BM haloes during the non-linear regime is followed by NB-TSPH simulations.

vi) The small cosmological box containing the perturbation eligible to become a galaxy is our approximation of the local Universe, the one subsequently affecting the evolution of the perturbation itself. The validity of this approximation will be examined in some detail below. To avoid numerical problems with sharp edges, we then single out the sphere of radius $r = l/2$ centered on the center of the box, change the particles coordinates from comoving to the proper physical values (this is simply achieved by dividing the comoving value by the initial cosmological expansion parameter $a = 1/\zeta_i - 1$, where $\zeta_i$ is the initial redshift provided by COSMICS at the end of the linear regime), and add a radial outward directed velocity component to each particle. This component of the velocity is proportional to the radial position of the particle and the initial redshift of the simulation. This velocity mimics the effects of the outward directed Hubble flow, to take the expansion of the Universe into account.

vii) To obtain the initial conditions for other model proto-haloes (galaxies) with different mass and/or mean initial densities, we start from the reference High Density, High Mass (HDHM) halo and proceed as follows:

Varying the mass. If the spatial dimensions of the box are changed, COSMICS modifies in a self-consistent fashion the size and consequently the mass of the haloes, leaving, how-

1 Alternatively, the equations of motion could be modified to account for the expansion. However, while this is straightforward to do in a comoving frame, it is less commonly done in physical coordinates. Moreover, virialized structures (like the model galaxies after the initial stages of evolution) are no longer subject to the expansion of the Universe and should be described by static coordinates. Examples of the method we adopt are given in [Katz & Gunn 1991, Kawata 1999] and in other similar studies.
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contrast \( \delta \) in the constrained field), because this would reduce the density of the constraint \textit{without} modifying the random perturbations, thus resulting in a very different global field with more spectral power on the small scales. To obtain intermediate density haloes, the mean density of the reference halo is decreased by a factor of 15%. The mean density of these latter is then decreased by another 15% to obtain low density haloes. Therefore, these latter have the initial mean density which is 72.25% of the value of the initial reference halo. Finally, a set of haloes with very low initial over-density was obtained reducing the mean density by 50%. As already pointed out, the different densities can be considered as a way to mimic the possible environments in which a proto-halo begins to form.

viii) At this stage, a minimal amount of solid-body rotation is also added to all the particles in the proto-haloes. The solid-body rotation is represented by the spin parameter

\[
\lambda = \frac{J/|E|^{1/2}}{GM^{5/2}}
\]

where \( J \) is the angular momentum, \( E \) is the initial binding energy, and \( M \) the total mass of the system. Typical values of \( \lambda \) range from 0.02 and 0.08 (White 1984), which corresponds to angular velocities of the order of fractions of a complete rotation over time-scales as long as about ten free-fall timescales (Carraro et al. 1998). We adopt \( \lambda = 0.02 \). The choice is motivated by the fact that our model galaxies are meant to represent ETGs. This issue will be also shortly commented in general remarks of this section.

ix) Finally, each particle in each proto-halo is split into a DM and a BM particle (gas), this latter with a small displacement from its original position to avoid numerical divergences. Indicating with \( m_0 \) the mass of the original particle and with \( f_{BD} = \rho_{BM}/\rho_{DM} \approx 0.1656 \) the cosmological ratio between the BM and DM densities, we assign to each gas particle the mass \( m_{gas} = f_{BD} \times m_0 \). Consequently, \( m_{DM} = (1 - 0.1656) \times m_0 \). The final HDHM proto-halo consists of \( \approx 58,000 \) DM particles plus an equal number of gas particles.

Coding the models. With this method, we have obtained twelve initial haloes. As the reference HDHM model, each one is identified by a string of four or five letters, of which the first two (or three) refer to its density and the last two to its mass (e.g., IDLM is for intermediate density, low mass; VLDHM stands for very low density, high mass, etc.). The key initial value of a few important parameters for all the twelve haloes are listed in Table 1 whereas a sketch of the density contrast inside the initial sphere projected on the XY plane for the model IDIM is shown in Fig. 1 where a number of lumps of matter are clearly evident. The number of particles in all models is similar to that of the HDHM case, with some minor differences due to the different displacement of the particles in the outskirts of the retailed sphere.

General remarks. We point out that all the proto-haloes we have chosen are fully consistent with the underlying cosmological background. The expected DM halo mass function follows the laws by, e.g., Press & Schechter (1974) or Sheth et al. (2001), and/or the analytical fits of numerical simulations (e.g., Warren et al. 2000). In this picture, a \( M \approx 10^{13} M_\odot \) halo is becoming “typical” on a \( \sim 10 \) Mpc
scale (the typical linear size of a large galaxy cluster) at redshifts below \( z \sim 5 \), which is compatible with the collapsing redshift of our HDHM model (see below). Such massive haloes are also expected not to be rare on ten times larger scales at \( z \sim 10 \). This can be easily checked considering the halo growth functions of DM haloes for the concordance Λ-CDM cosmology: the expected number of haloes with mass \( \sim 10^{13} M_\odot \) within a volume of \( 1 \) (Mpc/h)^3 at \( z \sim 10 \) is \( n \sim 10^{-8} \) (see Fig.2 of Lukic et al. 2007).

The main criticism that could be made to the above procedure is that the size of the simulation backing the perturbation in question is too small (9.3 comoving Mpc on a side for the HDHM box, and consequently smaller in the other cases). It is long known that the simulation size determines the maximum perturbation wavelength. If the long wavelengths are dropped out, the strength of subsequent clustering is reduced, but at the same time the number density of intermediate mass haloes (with total mass of the order of \( 10^{13} h^{-1} M_\odot \) in our case) is enhanced (Power & Knebe 2006). While this may be a problem in real cosmological simulations, in our case it is not so, because we are interested in objects with the mass and size of galaxies and not of galaxy clusters. Furthermore, Power & Knebe (2006) show that truncation of the initial power spectrum (i.e., the small size) of the simulation has little impact on the internal properties of the haloes. However, truncation lead to spin parameters that are 15\% lower than usual. We have in a sense avoided the whole spin problem just by taking the lowest value for the spin parameter (\( \lambda = 0.02 \)) suggested by White (1984), a reasonable approximation for slowly rotating systems such as ETGs. A final remark could be made related to the fact that adopting the sphere of radius \( r = l/2 \) (about the size of the Local Group) as the starting protohalo we implicitly neglect the possibility that late infall of nearby haloes that initially were outside the sphere may occur. This would inhibit late refuelling of gas (and stars and DM) to our system that is equivalent to say that we inhibit late mergers likely accompanied by star formation. This indeed is not the aim of the present study, which intends to explore the modality of star formation in alternative to the very popular hierarchical scheme. To conclude, the scheme we propose to derive the initial conditions for our model galaxies is not in conflict with the cosmological paradigm and it is easy to implement in NB-TSPH simulations.

2.2 The NB-TSPH code EvoL

EvoL is a Lagrangian N-Body parallel code designed to study the evolution of astrophysical systems on any spatial scale, from large cosmological volumes to sub-galactic regions. It is based on the standard Tree Algorithm for the description of the gravitational interactions Barnes & Hut (1986), and on a modern version of the Smoothed Particle Hydrodynamics method to model gas dynamics. Its basic features concerning the treatment of gravitational interaction and hydrodynamics are described in details in Merlin (2009) and Merlin et al. 2010; here, we briefly recall some of its characteristics.

EvoL includes some peculiar features, such as the lagrangian formulation of the SPH method that includes the so-called \( \nabla h \) terms in the equations of motion (taking into account variation of the smoothing length, see Merlin et al. 2010, for the definition), an artificial thermal conduction to smooth out contact surface discontinuities, and finally the variable adaptive softening lengths with consistent correcting terms in the equations of motion. In the present models we make the following assumptions and/or simplifications:

(i) We switch off the artificial thermal conduction, because its effects in combination with radiative cooling are not yet fully understood and tested, and may lead to unphysical behaviors, with particles exchanging heat because of numerical conduction and simultaneously losing energy due to physical cooling. The issue deserves an accurate analysis to be planned for a forthcoming study.

(ii) We adopt variable softening lengths; this is an important difference with respect to other similar calculations, in which a constant softening length has been adopted. The only exception is Nelson (2006) even if in that case no correcting terms in the equations of motion were considered. In the present models, we let the softening, \( \epsilon \), and smoothing, \( h \), lengths vary freely. In a few companion models calculated for the sake of comparison, we adopt for the two parameters the minimum threshold values \( \epsilon_{\text{min}} = 100 \) and \( h_{\text{min}} = 100 \) pc. Such low limits would provide sufficient resolution in all the cases of interest. However, these side models show that the results are very sensitive to these parameters, even if all the models are essentially similar in their global aspects (see Sect. 3.5 for a discussion on this issue).

(iii) We include a density-dependent pressure limit to avoid artificial clumping of poorly resolved gas. The method is similar to the one presented by Robertson & Kravtsov (2008). In practice, at each time-step all SPH particles are used to compute the local Jeans mass, which is a function of density and temperature, i.e.

\[
m_{\text{Jeans}} = \frac{4\pi}{3} \rho_{\text{gas}} \left( c_s^2 \frac{\pi}{G \rho_{\text{tot}}} \right)^\frac{3}{2},
\]

where \( c_s = \sqrt{\frac{\gamma}{\gamma-1}} u \) is the local sound speed, \( \rho_{\text{gas}} \) is the density of the SPH particle, and \( \rho_{\text{tot}} \) is the total local density (i.e., gas plus DM and/or stars). Then, the internal energy of each particle (limited to the computation of the hydrodynamical forces) is chosen as the maximum between the real value and the “effective” value, which is defined as the energy that the gas sphere should have to be stabilized against collapse:

\[
u_{\text{eff}} = N_{\text{SPH}} \times \frac{1}{7(\gamma - 1)} \rho_{\text{tot}} \left( \frac{3 N_{\text{SPH}} m}{4\pi^2 \rho_{\text{gas}}} \right)^\frac{\gamma}{\gamma - 1},
\]

where \( m \) is the mass of the particle, \( N_{\text{SPH}} \) is the number of particles required to resolve a region (indicatively, this number can be considered equal to the typical number of neighbours, i.e. about 60 in our models, or a multiple of this value), \( \gamma \) is the adiabatic index, and \( N_{\text{SPH}} \) is a free parameter \( \geq 1 \). In these models we use \( N_{\text{SPH}} = 15 \) (Merlin 2009; Merlin et al. 2012 in preparation). Note that \( u_{\text{eff}} \) is only used to compute the mechanical acceleration due to the internal gas pressure.

\[2\] This expression for the Jeans mass of a gas sphere in presence of a non collisional component can be obtained from a perturbative analysis of the stability of a pressurized cloud, see e.g. Tittley & Couchman (2000) for reference.
In the expressions for the variation of internal energy, dissipation, and cooling, the real value of $u$ is adopted. In this way, the temperature of gas particles can reach low values, without reaching high densities, thus creating regions of cold material without spurious numerical clumping of particles.

(iv) We run the simulations in physical coordinates since at present we are not interested in large scale motions. Moreover, this choice rules out the need for periodic boundary conditions, which would require a large amount of additional CPU-time.

(v) Our initial system is confined within the initial sphere of $l/2$ radius and it is surrounded by empty space. In other words the evolution of our model galaxy occurs in void. This simplification of the problem has some non-negligible drawbacks. First of all, the infall of material towards to innermost regions of the galaxy sooner or later stops because material outside the sphere cannot flow in. This rules out the possibility of late mergers and/or late infall of significant amounts of gas. Important point to keep in mind, but which does not invalidate the results of this study as long as late merger, matter acquisition are not of interest. Secondly, the feedback-heated gas leaving the galaxy will escape freely without experiencing drag forces. Thirdly, the density of the system cannot be consistently computed and the models evolve as small closed Universes rather than over-dense regions within a flat Universe, so the redshifts are only indicative of the real evolutionary time-line.

(vi) Finally, we include radiative cooling, star formation, and stellar feedback. As they are described in detail in Merlin (2009), in the sections below we limit ourselves to summarize their key aspects.

### 2.3 Radiative cooling

The cooling functions for atomic radiative processes are those described in Carraro et al. (1998), which in turn are based on those elaborated by Chiosi et al. (1998). In brief, for temperatures greater than $10^4$ K they lean on the Sutherland & Dopita (1993) tabulations for a plasma under equilibrium conditions and metal abundances $log[Z/Z_\odot] = -10$ (no metals), $-3$, $-2$, $-1.5$, $-1$, $-0.5$, $0$ (solar) and $0.5$. The chemical properties of gas particles are followed in details, tracking the abundances of Fe, O, Mg, Si and C, and the global metallicity Z. For temperatures in the range $100 < T < 10^4$ the dominant source of cooling is the $H_2$ molecule becoming rotationally and/or vibrationally excited through a collision with an H atom or another $H_2$ molecule and decaying through radiative emission; the data in use have been derived from the analytical expressions of Hollenbach & McKee (1979) and Tegmark et al. (1997), amalgamated together by Chiosi et al. (1998). We point out that in the present simulations it is assumed that all hydrogen becomes molecular below $10^5$ K, so the cooling rate in this range of temperatures may be somewhat overestimated. However, cooling from collisions of neutral hydrogen atoms with electrons from ionized metals (when present) is neglected. This may somewhat compensate for the excess of cooling due to molecular hydrogen. In any case this is a (marginally) weak point of our current treatment of cooling that should be improved.

Finally, for temperatures lower than $100$ K, Chiosi et al. (1998) and Carraro et al. (1998), starting from the studies of Caimmi & Secchi (1986) and Theis et al. (1992), incorporate the results of Hollenbach & McKee (1979) and Hollenbach (1988) for CO molecule as the dominant coolant. The following analytical relation in which the mean fractional abundance of CO is given as a function of $\text{[Fe/H]}$, is found to fairly represent the normalized cooling rate (i.e. $\Lambda_{\text{CO}}/n^2$ with $n$ the number density of particles)

$$\frac{\Lambda_{\text{CO}}}{n^2} = 1.6 \times 10^{-29} (\text{Fe/H})^{-1.699} T^{0.5}$$

in erg/s/cm$^3$. For all other details see Carraro et al. (1998).

We also include the contribution by inverse Compton cooling

$$\Lambda_{\text{Compton}} = 5.41 \times 10^{-36} n_e (T - T_{\text{CMB}}(1+z))(1+z)^4$$

in erg/s/cm$^3$, with the usual meaning of symbols (see e.g. Ikeuchi & Ostriker 1986).

The radiative cooling is computed separately from the SPH equation of energy conservation, due to its very short time-scales. During a dynamical time-step, the density and the metallicity of gas particles are kept fixed, as well as the mechanical heating rate. On the contrary, the temperature is let vary under the action of cooling, and the new cooling rates are simultaneously obtained as a function of the new temperature. The process is iterated until the end of the time-step.
2.4 Star formation

Star formation is modeled by means of a stochastic method, similar to that introduced by Churches et al. (2001) and Lia et al. (2002). First, only gas particles belonging to convergent flows (i.e. \( \nabla \cdot \mathbf{v} < 0 \)) and denser than a suitable threshold \( \rho_* \) are considered eligible to form stars. In our models we adopt \( \rho_* = 5 \times 10^{-25} \text{ g/cm}^3 \). No restriction on the temperature is imposed; this is motivated by the fact that thermal instabilities can produce star forming sites even within high temperature gas.

If a gas particle satisfies these criteria, it is assumed to form stars at the rate \( \frac{dp_i}{dt} = \epsilon_{SF} \rho_i / t_{ff} \), where \( t_{ff} \approx 0.5 / \sqrt{G \rho_{tot}} \) is the free-fall time, \( \rho_{tot} \) is the local total mass density (DM plus BM), and \( \epsilon_{SF} \) the dimensionless efficiency of the star formation process. This means that a gas particle is expected to transform a fraction \( \epsilon_{SF} \) of its mass into stars over its free-fall time scale. However, a stochastic description of the star forming process is adopted to avoid the creation of exceedingly large numbers of star particles. Thus, gas particles undergo a Monte Carlo selection to check whether or not they will actually form stars, in such a case they are instantaneously turned into collision-less star particles (see also Lia et al. 2002, Merlin 2009). To this aim, a random number \( r \in [0,1] \) is drawn and compared to the probability that the particle is actually forming stars. An obvious choice would be to use the probability \( P = \epsilon_{SF} / t_{ff} \times \Delta t \), where \( \Delta t \) is the dynamical time-step. However, the evaluation of the real probability must take into account that the dynamical time-steps are generally much shorter than the free-fall times of particles. This implies that within a free-fall time scale a number of star forming events and corresponding random draws are possible depending on \( \Delta t \), and therefore so does the global probability for a single gas particle to be turned into stars. If the global probability over \( t_{ff} \) is \( P \), the stochastic process of random selection must be corrected by considering that the probability summed over the number of draws \( n \approx \Delta t / t_{ff} \) must be equal to the probability of a single draw within the whole free-fall time. First we take as the probability within a single time-step the value \( p \approx P/n \), then we consider as the probability of a successful event over \( n \) draws the quantity \( p \times \sum_{i=1}^{n}(1-p)^{i-1} < P \). The correcting multiplicative term is

\[
 f = \frac{P}{n} \times \sum_{i=1}^{n}(1-\frac{f}{n})^{i-1} \tag{7}
\]

Therefore, the final probability during the time-step is \( f \times p \). If \( r \leq f_p \), then the gas particle is instantaneously turned into a stellar particle.

In the present models, we assume \( \epsilon_{SF} = 1 \). Empirical estimates of the efficiency of star formation based on observational data of star forming events inside molecular clouds in the local vicinity yield \( \epsilon_{SF} \approx 0.025 \) (Lada & Lada 2003; Krumholz & Tan 2004). In this case the gas density and free-fall time scale of are those of the cold molecular clouds. However, in the case of a large scale star formation mechanism in a galaxy, the above estimate may not correspond to reality. As a matter of facts, considering that a typical galaxy with \( 10^{11} M_\odot \) mass in stars has to disposal a time scale of about 13 Gyr to build up its stellar content, the mean estimate of the star formation efficiency is closer to 0.1 rather than 0.02. The new estimate would increase by nearly a factor of ten if the time to disposal to form stars is much shorter than 13 Gyr, say 1 to 2 Gyr. In addition to this, it is not known whether \( \epsilon_{SF} \) should be maintained constant during different cosmic epochs due to the different mechanisms forming stars at high redshifts and in metal poor regions with respect to local molecular clouds. Finally, it is not known whether the efficiency of star formation is the same in all galaxies independently of their mass (either total or baryonic). For all these reasons, we prefer to adopt here \( \epsilon_{SF} = 1 \), which means that the process of star formation occurs at 100% efficiency (meaning that all gaseous particles satisfying the SF criteria will turn their total mass into stars within a free-fall time scale). Another consideration has to be made here. Several numerical simulations calculated with different values of \( \epsilon_{SF} \) clarify that within a certain range of values, the star formation histories of the models with small values of \( \epsilon_{SF} \) are apparently quite similar to those with \( \epsilon_{SF} = 1 \), the only difference being that with small values of \( \epsilon_{SF} \) the simulations become much more time-consuming. The reason is that dense and cold clumps of matter, interacting with hotter material heated up by close SN explosions, require extremely small time-steps. In contrast, if the gas can easily form stars, this critical situation is avoided, and the simulations proceed much faster. The weak impact of \( \epsilon_{SF} \) on the final SF history can be attributed to the self-regulation cycle between star formation and energy feedback. A high SF efficiency implies a strong and sudden burst of stellar activity; but young stars soon pressurize their surroundings via energy feedback (see below), quenching the formation of new stars. However, if the gas is sufficiently dense, radiative cooling is effective and further star formation can soon take place (positive feedback). With a low efficiency, less stars form. Their heating is consequently lower, and more stars can soon form before the feedback halts their formation. This ultimately leads to the same situation as in the previous case, with perhaps the minor consequence of a delayed enrichment in heavy elements of the medium. Clearly, other parameters play a more important role; for example, the density threshold \( \rho_* \) and the efficiency of feedback. Therefore the choice \( \epsilon_{SF} = 1 \) and fast calculations likely, is the best compromise. However, one should always keep in mind that adopting \( \epsilon_{SF} = 1 \) may have other consequences on the dynamical evolution of the systems, favouring the collision-less collapse instead of the dissipative one (see Sect. 3).

2.5 Energy feedback

When a gas particle is turned into a star particle, it can be considered to represent a Single Stellar Population (SSP) made of many real stars. It starts re-fueling the ISM with heavy chemical elements and energy, mainly because of winds from young massive stars and Supernova (SN) explosions.

Winds. We consider two regimes for the energy injection by winds. When the SSP is young, the main source of energy are the young massive stars. Their winds have very high speeds, typically from 1000 to 3000 km/s. Here we adopt a constant velocity of \( v_{YSO} = 1500 \text{ km/s} \). The kinetic energy of the winds is assumed to be thermalized and released within the surrounding medium with an effi-
ciency of 20% (see e.g. Dyson & Williams [1997]). In the late stages of the SSP evolution, slow velocity winds from low mass stars become predominant; in these cases we assume the wind speed of $v_{OSO} = 60$ km/s (typical of RGB stars), and the same thermalization efficiency.

**SN explosions.** Each real SN explosion is expected to deposit some $10^{51}$ kinetic ergs in a very small region, and on a short time scale. However, most of this energy is soon radiated away, and only a small fraction of it is subsequently thermalized. We use the results by Thornton et al. (1998) to obtain an analytic approximation of the fraction of the initial energy which becomes available at the end of the expanding phase of the SN bubble. However, in the real Universe SN explosions take place in already shock-heated regions, with temperatures raised up to some $10^6$ K, because of the action of a photo-ionizing flux from massive young stars, stellar winds, and previous SN explosions. All this is not taken into consideration in Thornton’s study, so a large uncertainty still affects the description of the whole process. However there are a few exceptions like Cho & Kang (2008) that should be taken into account. A single star particle represents an entire SSP, so a large number of SN explosions are expected to take place within a single star particle over a rather long time scale. The number of SN explosions (attention must be paid to release the energy in discrete bursts consistently with the number of real explosions), as well as the amount of gas released by a SSP of given age and mass during a time-step and its chemical composition are computed with the technique described in Lia et al. (2002), using theoretical SSPs (in EvOL, the Padova tracks are adopted, see Greggio & Renzini [1983]), and adopting the initial mass function for the real stars in the SSP (star particle) of Kroupa (2001). We refer the reader to the above papers for all the details. It is worth clarifying that we do not adopt stochastic approach of Lia et al. (2002) to model SN explosions. At each time-step they decided the fate of each star particle by means of the Monte-Carlo method; randomly chosen star particles were instantaneously turned back into gas, while the remaining ones continued their life as star particles. On the contrary, in our description each SSP continuously releases gas and energy at each time step, becoming a hybrid particle (in the sense described below).

Finally, the total budget of thermal energy produced by stellar winds and SN explosions is given continuously to a sort of gas reservoir assigned to star particles, ultimately formed by the gaseous material ejected from the SSP. This reservoir, due to its high temperature, acts a piston on the neighboring particles, effectively injecting kinetic energy into the surrounding interstellar medium. In the next Section we clarify some aspects of this method.

### 2.6 Fate of the gas ejected by SSPs

The gas ejected by SSPs, heated up by the energy budget from winds and SN explosions and enriched in chemical elements, interacts with the surrounding medium. This is a crucial issue in numerical simulations of galaxy formation. Since it is not feasible to continuously form new gas particles from the stellar particles, it is necessary to develop a consistent, yet simple method to model this process.

There are two main sources of uncertainty. First, one has to decide how to deal with the hot gas *soon after* its ejection from the SSPs (this gas is essentially composed by matter ejected by stars in form of winds and/or SN explosions). How should it be let cool? How to compute its density? It has long been known that distributing the feedback energy among neighboring particles and then letting them normally evolve among the other cold ones yields poor results and very low feedback efficiencies [Katz 1992]: the energy is soon radiated away and, in critical situations, hot particles escape from the region due to excessive hydrodynamical friction with their neighbors that is not physically grounded. Also, distributing the gas released by the SSPs to nearby SPH particles at each time-step has proven to be inefficient, since the large amounts of energy released locally are smoothly spread out too strongly. To cope with all this, we model we proceed as follows. We suppose that initially the ejected gas remains stuck to the parent star particles. This assumption sounds reasonable in very early stages of gas ejection, but surely becomes unphysical as time goes by. Sooner or later, the gas ejected by SSPs should move freely and mix with its nearby particles. Therefore, the issue is how and when the gas should be let evolve “normally”, as a standard SPH gas particle - i.e., when it should dynamically decouple from its parent star particle.

To this aim, we split the evolution of the ejected gas into two temporal steps. The gas is kept locked to the parent SSP during the first stages of its evolution, namely from its first appearance in the SSP to the age of $t_s \approx 8 \times 10^7$ years. This is the typical lifetime of a $6 M_\odot$ star with solar composition and convective overshooting ([Bertelli et al. 1994]). In this type of stellar models, the $6 M_\odot$ star is also the lower limit for the occurrence of Type II Supernovae. The value of $t_s$ could be lowered by a factor of 2 (and the mass limit accordingly increased to about $8 - 9 M_\odot$) to be more conservative, however without changing the overall picture we are proposing. During this first stage, the ejected gas does not contribute to the acceleration of the parent

![Figure 2](image-url)
heavy elements (namely, Fe, Mg, O, Si, C, and global metallicity Z) ejected by each SSP, i.e. star particle, is computed at each time-step by means of the Padova evolutionary tracks and the Greggio & Renzini (1983) rates based on the number of SN events and the amount of metals released by the “dying” SSPs during the current time-step. The global chemical yield from a single SSP during each time-step is given by the elements released by dying (exploding) stars during that time-step. These are the sum of three contributions: (i) the elements locked in the dying stars at the moment of their birth, inherited from the proto-stellar gas; (ii) the heavy elements created in the core by thermonuclear reactions and ejected in Supernova II explosions; (iii) the elements created in the core of binary stars which end their life as Supernovae Ia. The reader should refer to Lia et al. (2002) for a detailed description of the method.

Numerically, the heavy elements released by a SSP are assigned to the gaseous mass ejected by the same SSP during the current time-step and subsequently spread over neighboring particles by means of a diffusion algorithm based on the Fick’s law in spherical symmetry:

$$\frac{∂f}{∂t} = D_{\text{diff}} \frac{∂^2 f}{∂r^2}.$$  (8)

The formalism is the same we have adopted for thermal conduction due to diffusion of electrons. The coefficient $D_{\text{diff}}$ is obtained from the diffusion velocity of electrons by Monaghan (1992).

2.8 Re-ionization and the first generation of stars

Cosmic re-ionization, as inferred from the WMAP-5 data, is currently believed to have taken place at $z \leq 10$. In our models, we include a crude treatment of photo-heating from re-ionized free electrons. Following Pawlik & Schaye (2009), we increase the temperature of all gas particles fueling their internal energy reservoir with an additional budget of 2 eV per proton, which is instantaneously injected at $z = 9$. The effect of this energy boost is to raise the temperature of the low density particles to some $\sim 10^5$ K. In high density regions, however, radiative cooling is so efficient to keep the temperature well below this value. Fig. 3 compares two snapshots taken from models in which this re-ionization effect is (is not) included. However, it is worth noting that this way of proceeding lets the gas particles free to cool down soon after the sudden injection of energy; consequently, the cos-
mic web may happen to get too cold at later times (this can be seen in Fig. 13 of Sect. 3.3).

In this context, we also include in a very simple fashion an early population of metal-free stars otherwise known as Population III (Pop III) stars (e.g., Bromm et al. 2002; Yoshida et al. 2003 and references therein for more details). We assume that when a metal-free gas particle is prone to form stars for the first time (see Sec. 2.4) it generates only Pop III objects. Furthermore, this SSP is supposed to contain only massive stars, because according to current belief the underlying initial mass function is heavily skewed towards massive stars (Bromm et al. 2002). Consequently the Pop III generation of stars is short lived and within a few million years (almost immediately on galactic time scales) injects lots of energy into the interstellar medium by stellar winds and supernova explosions. Therefore, thanks to this we can simplify the occurrence of this first generation as follows: gas particles undergoing star formation are not turned into stars, but simply raised to high temperature, $T \simeq 2 \times 10^7$ K, and enriched in metals, $Z \simeq 10^{-4}$. By doing this we do not care to track the fate of Pop III stars as far as their remnants (ultra-compact objects and Black Holes) are concerned. Since these remnants could be the seeds to form to a central super-massive black hole, most likely fueling a period of AGN activity, a complete description of the remnants from Pop III stars is in progress.

2.9 General remarks

Before going into the results of our study, we shortly discuss here a few possible drawbacks of our models.

First, as already mentioned, we adopt the stellar efficiency $\epsilon_{SF} = 1$, in contrast to what observational data and simple theoretical considerations would suggest. This has proven to cause non-negligible effects on the dynamical evolution of the systems; they will be discussed in the Sections below.

Second, no treatment of the multi-phase nature of the interstellar medium has been included. In contrast, different phases naturally develop within the gas (see Sect. 3.1). Furthermore, it is long known that numerical effects like overcooling may arise if the gas is modeled according to the standard SPH method (Merlin & Chiosi 2007). Given these considerations, a more detailed treatment of the interstellar medium not only is urgent but may also cure part of the problems encountered with overcooling.

Third, we have not considered the presence of nuclear activity, neglecting both the dynamical action of a central super-massive black hole, and the energy feedback from an AGN, both of which are currently believed to play important roles in the evolution of a proto-galaxy.

Fourth, we have not included other sources of feedback apart from stellar - i.e., magnetic fields and their pressure or cosmic rays.

Finally, our cooling functions do not take into account the state of the Hydrogen gas; in particular, the variation of the molecular fraction of Hydrogen, which strongly influences the cooling efficiency below $10^4$ K in primordial, non-enriched gas (see Sect. 2.4).

3 MODEL RESULTS

All the simulations have been run on the HP Blade Server MONSTER with 14 hosts and 116 CPUs at the Astronomy Department of Padova. Each major galaxy model made use of four parallel CPUs, whereas companion test models of small size were run on single processors. On the whole, the simulations took $\sim 100,000$ CPU hours, requiring from 30,000 to 80,000 time-steps. There are two groups of models:

First group: testing the models. The first group contains the test computations performed to assess the model response to varying some important parameters. They are calculated with a smaller number of particles (usually from 5000 to 10000). To save computational time, not all the models have been calculated up to the present age. However all of them are carried out to the stage in which the mass assembly is completed (say at least a few Gyrs). The following questions are addressed:

(i) Star formations efficiency. These are the models already described in sect. 2.4. They are calculated to check the sensitivity and overall behaviour at varying the SF dimensionless efficiency $\epsilon_{SF}$ of the star formation rate.

(ii) Stellar feedback. A model is run with no feedback from stars. As expected, the star formation rate grows dramatically at early times (see Fig. 4; the run had to be stopped very soon because the continuous collapse of cold gas, not contrasted by the action of heating due to feedback, requires smaller and smaller time-steps and therefore high computational costs). Our prescription for the stellar feedback can therefore be considered as reasonably efficient.

Second Group: the reference models. The second group contains the reference set of models, with higher resolution. All of them are calculated well beyond the completion of the mass assembly, a few even down to the present age or very close to it. All important data referring to the models at the same reference redshift $z = 1$ are summarized in Table 2 whereas in Table 3 are displayed similar data but for the last computed stage of each model (different age and redshift).

Results for the reference models. In the following we shortly describe the main results for the twelve reference models. We start showing in Fig. 5 a snapshot of their geometrical structure projected on the $XY$ plane at the last computed. In most cases the spatial distribution of stars resembles that of classical ETGs; in one case a filamentary substructure, and perhaps satellites companions are visible. Fig. 6 displays the star formation histories of the twelve models along their lifetime, down to the last simulated epoch. As found by Chiosi & Carraro (2002) and expected here, the kind of star formation at work changes with the galaxy mass and initial density. It is monolithic-like (a single dominant initial episode) in high mass and/or high initial density proto-galaxies, it gradually turns into a broad rather long and mild activity at decreasing mass and/or initial density, and finally it gets like a series of discontinuous very episodes (bursting mode) for low mass and/or low initial densities galaxies. The point will be developed in more detail below.

3.1 Dynamical evolution and final morphologies

All the models go through an initial phase of expansion, as they follow the Hubble flow. The central regions, in which
Figure 5. The spatial distribution of the stellar content of the model galaxies at the redshift of the last computed stage, projected on the XY plane. Left to right, top to bottom: HDHM ($z = 0.22$), IDHM ($z = 0.77$), LDHM ($z = 0.49$), VLDHM ($z = 0.83$); HDIM ($z = 1.0$), IDIM ($z = 0.75$), LDIM ($z = 0.58$), VLDIM ($z = 0.15$); HDLM ($z = 0.36$), IDLM ($z = 0.22$), LDLM ($z = 0.05$), VLDLM ($z = 0.0$). The coordinates X and Y are in proper Mpc.

the density peak is more pronounced, soon detach from the outward motion and start re-collapsing, forming the core of the virialized structure. As time goes on, more and more external regions turn around and fall onto the central virialized core. The process can be roughly described as a secondary infall, with time-scales strongly dependent on the properties of the particular halo under consideration. More massive haloes globally virialize later with respect to smaller ones, as expected in the hierarchical cosmology. However, denser haloes also re-collapse later than the shallower ones, so the epoch of virialization is decided by the interplay between these two effects. At times depending on the depth of the central potential well, the central regions of the haloes begin to reach the critical density $\rho_\star$ at which the gas becomes able to form stars. Due to the strongly peaked density perturbations, all models soon develop a central stellar system; however, in the very first stages of the evolution we note the formation of many small clumps of stars, which soon merge into a single entity. In most cases, additional smaller stellar systems form in regions far from the center because of local concentrations of cold gas, eventually merging onto the major body of the proto-galaxy. In model galaxies of low mass, sparse isolated stars form within clumps of expanding gas (wind), forming extended haloes of extremely low density around the central body of the proto-galaxy.

The whole history of the building up of the stellar component of the model galaxies is shown in series of panels in Fig. 6. Each panel show the projection on the XY plane (coordinates in proper Mpc) of the position of the star particles. Each row corresponds to a total mass (high, intermediate and low starting from) and each columns to an initial over-density: from high (left) to very low (right). Finally each group contains four subpanels depending on the redshift at which the snapshot are taken: from left to right and top to bottom the redshift is $z = 11.5$ (7 for the very low density, high mass case), 5, 3, 1.

All models finally develop a central dense stellar system, located at the centre of the collapsed halo of DM. Some amount of spurious global bulk motion of the system (due to numerical instabilities) can be observed in some cases (see Figs. 5 and 6). At $z = 1$, most of the systems have almost completely relaxed into stable configurations, although in the massive ones pronounced irregular features are clearly visible (arms of star forming regions, diffuse haloes of stars around the central object). These are not formed by old populations infalling onto the central object, but by young SSPs escaping from the galaxy (this can be inferred by checking the birth age of the star particles and their velocities). The bright, spherical stellar clump visible in the bottom of the VLDHM model in Fig. 5 has mass $M_\star \approx 10^{10} M_\odot$, much larger than the masses of the low mass models. Such stellar clumps must therefore be considered tidal galaxies, formed by gas condensation. Many examples of such systems are known to exist (see e.g. Elmegreen et al. 2007), even if they are generally thought to be generated within interacting systems. In our cases, local condensations of gas, compressed...
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Figure 6. Star formation histories for the twelve reference models. Left to right: high density, intermediate density, low density, very low density. Top row: high mass; intermediate row: intermediate mass; bottom row: low mass.

and pushed outwards by the pressure exerted by the inner regions of the system, collapse to form stars while being ejected from the galaxy. The tidal objects are finally engulfed by the massive galaxy, but usually leave a faint expanding trail of debris.

On the other hand, low mass systems are surrounded by extended haloes of sparse stellar particles, as shown in Fig. 8 for model HDLM at \( z = 1 \). Perhaps, these may be identified with the observed intra-cluster light (ICL, Monaco et al. 2006), which could therefore be the result of diffuse star formation within gas escaping from low mass galaxies, rather than the consequence of tidal stripping and/or disruption of bounded systems.

It is worth recalling here that a different choice for the dimensionless efficiency \( \epsilon_{SF} \) may have lead to different dynamical histories, as discussed below. In particular, a lower efficiency could have reduced the formation of diffuse stellar haloes.

All galaxies have triaxial shapes; their axis ratio at \( z = 1 \) and at the final redshifts, obtained from their projection on the \( XY \) plane, are summarized in Tables 2 and 8. If observed from this angle of view, all the models would be classified as E3 - E5 galaxies in the Hubble system. Obviously, more spherical shapes could be obtained changing the view angle, and almost all models could be morphologically classified as almost perfect spherical galaxies choosing a suitable projection. Noticeably, the minimal initial rigid rotation can hardly be considered the main cause for the pronounced triaxiality of the models, which must be therefore a consequence of their dynamical history. In particular, observing the first stages of the formation of these systems it appears that their collapsing motions took place along the filamentary structures of the cosmological over-densities, and stars tend to conserve elongated orbits. This view is strengthened by the fact that reducing \( \epsilon_{SF} \), by requiring a higher degree of dissipation and condensation of gaseous particles, yields more concentrated and spherically-shaped structures. It would be worth investigating how this would affect to present day structure these systems. Perhaps, in this context, a non-negligible role could also be played by the absence of large scale tidal motions caused by the adopted prescription for the evolution of the whole proto-galaxy (the initial sphere) in void.

The injection of energy from stars into the ISM causes turbulence and large scale motions in the gas (see for instance the situation illustrated in Fig. 2). Depending on the particular circumstances under which it takes place, the stellar energy feedback can lead to the ejection of large amounts of gaseous material from the potential well of the galaxy, or to the formation of condensed structures, resulting in new star forming sites. The minor impact of feedback in low mass haloes ultimately results in the large differences in their star formation histories, as explained below.

Different phases naturally develop in the gas during the first stages of the galaxy formation as clearly shown the phase diagram (temperature vs. density of the gas) displayed in Fig. 9 for the prototype model IDIM. First, the initially cold, expanding gas begins to heat to a typical temperature of \( T \approx 10^4 \) K in the central regions, as soon as they detach from the Hubble flow and collapse. When the radia-
tractive cooling becomes effective, this gas begins to cool and to form stars. The ejecta from the SSP subsequently heat the surroundings\(^3\). Now, the infalling gas is forced to interact with these heated regions. If the hot gas is enough, this may lead to expanding shock fronts (see below). Thus, after several Myr of activity, the gas can be roughly split as follows. A cold and dense phase is formed by gas in the central regions of the galaxies, where stars are forming. On the opposite extreme, hot and thin gas \textit{either} is leaving the halo (galactic wind) after being heated by stellar feedback, \textit{or} it is infalling onto the central object, after being shock-heated because of the interaction with the central, hot regions. Two more phases are clearly detectable: the hot \textit{and} dense gas, where the stellar feedback is heating the gas at very high temperatures before these regions expand under the action of the local pressure, and the cold but thin gas belonging to the cosmological cosmic web (before re-ionization takes place, or cooling after re-ionization; see Sect. 2.8).

3.2 Star formation histories

Looking at Fig. 6, it is clear that the initial conditions of each halo deeply influence the star formation history (SFH) of the parent proto-galaxy. Even at a first glance, two clear trends can be easily identified.

First: the more massive the halo, the more peaked is the SFH. All massive haloes \(M_{\text{tot}} \simeq 10^{14} M_\odot\) show a single, strong burst at very early times, followed by a slow decrease down to quiescence. A weak tail of activity persists down to later times in the less dense systems, thus accounting for a weak dependence on its density (or environment). Smaller haloes, in turn, display a more fragmented and prolonged
star formation activity, often alternated with periods of quiescence.

Second: lower densities correspond to delayed activity (the first burst starts later) and minor overall efficiency (the peaks of activity have lower magnitude). A delay up to $\approx 1$ Gyr in the beginning of the SF activity can be seen in the low density models. Moreover, the maximum rate is generally decreased and the activity is prolonged (this second effect is of crucial importance in intermediate mass haloes $M_{\text{tot}} \approx 10^{11} M_\odot$). Apparently, the mass of the halo is the dominant
Figure 9. Phase diagram density $\rho$ (g/cm$^3$) vs. temperature T (K) of gas particles for the IDIM model, at four stages of evolution ($z \approx 10.0, 7.0, 5.0, 3.5$). The hot gas decreasing its density is leaving the galaxy, due to stellar feedback.

Figure 10. Star formation histories of the whole galaxies (solid black line, large circles) compared with the corresponding ones within the central region ($r_{\text{inner}} \approx 20$ kpc; red dashed lines, large points) for two large mass systems. Left panel: model HDHM. Right panel: model VLDHM.

characteristic to determine its evolution, but intermediate mass haloes can have very different histories depending on their global over-density.

Table 4 summarizes the properties of the SFHs. Interesting quantities are the intensity of star formation (in $M_\odot$/yr) at its peak value and the time (in Gyr). Furthermore, the gradual building up of the stellar mass in a galaxy is measured by the percentage $p_*$ of the total stellar mass at redshift $z=1$ assembled at selected values of the redshift ($z\approx10, 5, 2$ and $1.5$), and the time (redshift) at which 50% and 99% of the stellar mass at $z=1$ is assembled. This is indicated by the pairs $(t, z)_{f=1/2}$ and $(t, z)_{f=99/100}$ in Table 4.

All the models begin forming Population II stars between $z \approx 20$ and $z \approx 7$. The HDHM halo experiences the strongest and earliest peak of activity, reaching some $1.5 \times 10^3 M_\odot$/yr at $z \approx 10$. More than 50% of its total stellar mass is assembled before $z \approx 6$. If observed now from the Earth, this galaxy would appear massive ($M^* \approx 3 \times 10^{11} M_\odot$) and essentially red at $z \approx 4$. Its tail of SF activity continues down to $z \approx 1.6$, contributing 19% of the total stellar mass. Looking at less dense haloes of the same mass, it can be noticed how the maximum of activity decreases in magnitude and is delayed to later times, down to a peak of $7.5 \times 10^2 M_\odot$/yr at $z \approx 5$ in the VLDHM model, where moreover an intense activity is prolonged to much later epochs. Despite the similarity between the final objects formed from massive haloes, a clear trend can be extrapolated from the data given in Table 4: the lesser dense the halo, the later the mass assembly takes place (with the VLDHM halo assembles half of its total stellar mass at $z \approx 2.7$). The overall efficiency of the star formation process in these models is 20%-30%.

Intermediate mass systems apparently have more complex histories. Moving from denser to less dense systems, the shape of the star formation curve changes dramatically. Strongly over-dense haloes undergo an intense burst at early times, much as the most massive systems. However,
less dense haloes gradually display more prolonged, and uneven histories, with many different peaks. Apparently, in this range of masses the initial density plays a dominant role in deciding the fate of the mass assembly of the galaxy. The efficiency of the star formation process is very high in these cases: 30\%-50\% of the initial gas is converted into stars.

All low mass systems show a continue, bursty SFH, sometimes with long periods of quiescence. In these systems, the feedback from stars is sufficient to halt the star formation process soon after its beginning. However, this means that as soon as the first stars cease their strong feedback activity, the gas is able to re-collapse and form new generations of stars. This galactic “breathing” leads to systems in which many different generations of stars are present. The overall efficiency of the star formation mechanism is lower in this cases, with 10\%-20\% of the initial gaseous mass turned into stars (giving high $M/L$ ratios, comparable to those of many observed dwarf galaxies).

Figs. 10 and 11 compare the SFH of four models (HDHM, VLDHM, HDLM, VLDLM) across the whole systems with the SFH within their central regions (that is, within the central 20 kpc in massive systems and within the central 2 kpc in the small ones). It can be noticed that massive systems build their central regions before their outskirts, while in less massive galaxies the SF process is active both in the inner and in the outer regions during the whole lifetime of the galaxies. This is consistent with some recent observational findings (van Dokkum et al. 2010) on the way galaxies are built-up.

3.3 Stellar populations: ages, mean metallicities, and gradients in metal content

Stellar ages. The distribution of the galactic ages ($T_{G,i}$) at which the star particles are created as a function of the radial distance is shown in Fig. 12 limited to the high and low mass models (the behaviour of the intermediate mass ones falls in between these two). The real star particles ages are $T_\ast = T_{G,i} - T_{G,i}$. In early epochs, the stars are preferentially created in the central regions, then the star forming activity expands to larger radii (inside-out mechanism), and moving towards the present time, the stellar activity tends to shrinks again towards the centre. This simply mirrors the star formation history and the mechanism of mass assembly presented above.

Mean metallicity. Turning now to the metal content...
reached by the model galaxies, we look at the average stellar metallicity as a function of the total stellar mass. The results are shown in Fig. 13, where the red circles refer to all the star particles, whereas the blue ones only to those within the central 5 kpc of each model. The black lines are the median (solid) and the 16th and 84th percentiles (dashed) of the data presented in (Gallazzi et al. 2005). There is marginal agreement between data and theoretical results. The absolute values of the theoretical metallicities are too low compared with the observational estimates, in particular for the most massive models which are almost an order of magnitude too metal-poor. This may be due to several factors. First, the theoretical values are straightforwardly computed averaging the metallicity of all particles within the chosen radius without any further refinement; of course, observed metallicity are luminosity weighted, while in our models the contributions from very metal-poor, primordial (old) and faint stars the same weight of more recent stellar populations. Second, part of the discrepancy may arise because of the different methods used to obtain observational and theoretical metallicities. For example, the use of the Kroupa IMF, which is tailored to fit the solar vicinity, may lead to underestimate the enrichment by massive stars in other environments. Recent studies seem to indicate that the IMF in massive ETGs is more skewed toward massive stars than commonly assumed (Cappellari et al. 2012) thus implying a net increase of metal production by massive stars in these systems. Third, the high star formation efficiency can again play a role, favouring the formation of low metallicity stars at early times and reducing the average metallicity of the models. Finally, as shown by Scodeggio (2001) and discussed by Tremonti et al. (2004), at least some part of the observed color- (and hence metallicity-) mass relation might just be a spurious consequence of an aperture effect, as observed colors (and metallicities) are usually measured within a fixed aperture, rather than a given fraction of the galaxy effective radius, so that the absolute values are not well determined. In conclusion, considering that chemical enrichment in NBTSPH simulation is still far from being fully satisfactory, we are inclined to say that theory and data marginally agree with each other.

In passing, we point out that the models naturally reproduce the observed trend of the α-enhancement phenomenon in massive galaxies. The early peak of star formation in high mass systems, followed by quiescence, ensure the enhancement in light elements, whereas low mass systems with prolonged SFH are less enhanced, as it should be (see the discussion in Tantalo & Chiòs 2002).

Metallicity gradients. Radial gradients in spectro-photometric properties of ETGs are known to exist (see e.g. La Barbera et al. 2010, and references therein). The majority of ETGs in the local Universe feature red cores (La Barbera & de Carvalho 2009), and less frequently blue cores, mostly found at low mass (Suh et al. 2010). The radial gradients are interpreted as age and metallicity gradients of the constituent stellar populations (e.g. González 1993).

In Fig. 13 we plot the metallicity of each particle as a function of projected distance to the galaxy center (from left to right, and top to bottom, models LDHM, HDHM, LDLM and HDLM are displayed). For each model, the median-binned metallicity profile is shown, the number of particles being the same in each bin. In order to estimate the metallicity gradient, we performed an orthogonal least-square fit of the metallicity profile in the range 0.1 r_e to 1 r_e, i.e. the same range used in observational studies (Poel et al. 1994). The fitting was repeated for 100 different two-dimensional projections, the average slope of the best-fitting lines giving the metallicity gradient, \nabla_Z = d \log Z / d \log R, of a given model. We repeated the computation of \nabla_Z by applying different weighting schemes, where each particle was assigned no-weight, mass-weighted, and luminosity-weighted, finding no significant differences within the errors, especially at high mass (see Tab. 5).

Looking at the results in some detail, high- (low-) mass models exhibit metallicity gradients spanning the range from \(-0.4 \pm 0.02\) to \(-0.27 \pm 0.16\), depending on the weighting scheme and environment. In general, low-density models tend to have more negative metallicity gradients. Moreover, at given density, high- (relative to low-) mass systems have more negative gradients.

In general, the majority of the stars in the central regions of the massive systems have higher metallicity than those in the outskirts; on the contrary, in less massive systems the metallicity of the central stars is generally equal to that of the peripheral ones. Even if they are not straightforwardly visible in Fig. 13 interesting features to recall are: (i) on the average, the metallicity distribution of the star particles gets narrower at increasing radial distance; (ii) the maximum values are always reached in the centre, (iii) great deals of stars are made of recycled gas; (iv) finally the simulations often show the presence tidal satellites which have their own chemical history.

Although the metallicities of our models do not match observations in a absolute sense, the relative variations of metallicity (between, e.g., different masses, and/or different galacto-centric distances) are likely more robust, as they are less dependent on the absolute calibration of the ingredients the models rely on. This motivates for a comparison of the predicted \nabla_Z’s with observational results. A steepening of the metallicity gradient with galaxy mass, qualitatively consistent with that seen in our models, has been reported by (La Barbera et al. 2010), and, independent of galaxy environment, by (La Barbera et al. 2011). On the other hand, other observational studies, based on smaller samples, do not find a clear correlation of \nabla_Z with mass (Spolaor et al. 2009), or even a double-value behavior (Tortora et al. 2010). For a galaxy mass of \(3 \times 10^{11} M_\odot\) (similar to that of \(7 \times 10^{11} M_\odot\) of models HDHM and LDHM), La Barbera and collaborators found metallicity gradients of \(-0.37 \pm 0.02\) and \(-0.41 \pm 0.02\) for low- and high-density ETGs, respectively. At low density, the estimate of is fully consistent with our model predictions (for no-weight and mass-weighted \nabla_Z’s). At high- (relative to low-) density, the models exhibit shallower gradients, by \(\sim 0.07\), in contrast to the data, where a (marginal) difference of \(-0.04 \pm 0.03\) is found. For what concerns low-mass models, they exhibit almost null gradients, consistent with the observational finding of Spolaor et al. 2009 that the \nabla_Z vanishes at a galaxy mass of \(\sim 10^9 M_\odot\). Whilst these results point to a fair agreement between our models and (some) available data-set, we postpone a more thorough investigation of both metallicity, age, and color gradients, as well as their correlation with other model properties, to a forthcoming contribution.
Table 5. Metallicity gradients of four ETG models in the radial range of 0.1 r_e to 1 r_e. Column 1 is the model ID label. Columns 2, 3, and 4, correspond to cases where the ∇Z is computed with no-weight, mass-weight, and luminosity-weight assigned to each particle. Errors are the rms of ∇Z estimates among 100 projections of each model.

| ID    | ∇Z (no-weight) | ∇Z (mass-weighted) | ∇Z (lum-weighted) |
|-------|----------------|---------------------|-------------------|
| LDHM  | −0.39 ± 0.04   | −0.40 ± 0.04        | −0.33 ± 0.03      |
| HDHM  | −0.32 ± 0.04   | −0.31 ± 0.04        | −0.27 ± 0.03      |
| LDLM  | 0.00 ± 0.04    | +0.01 ± 0.02        | 0.00 ± 0.06       |
| HDLM  | +0.09 ± 0.02   | −0.02 ± 0.01        | 0.16 ± 0.03       |

3.4 Cold and hot gas accretion and ejection (galactic winds)

Fig. 15 shows the gas temperature vs. density diagram for models of our set at z = 2. As expected from theoretical considerations (Rees & Ostriker 1977), a bimodal gas accretion can be observed in our models, as a function of the total mass of the proto-galactic halo. In low mass haloes, the gas accretes from the cosmic web without being shock heated to the virial temperature, and proceeds to flow along filaments towards the center of the halo, where it will eventually shock. In contrast, massive haloes soon shock the infalling gas to very high temperatures, suppressing the formation of local substructures and forming a central distribution of hot gas. haloes which do support shocks close to the virial radius are expected to contain a quasi-hydrostatic atmosphere of hot gas. Birnboim & Dekel (2003) claim that shocks can only form close to the virial radius in haloes with mass greater than 10^{12} M_⊙ for primordial gas (or around 10^{12} M_⊙ for gas of Solar metallicity), in good agreement with our models.

Looking at Fig. 15 it can be noticed that the situation shows a strong dependence on the mass of the halo. In massive and dense systems, almost no cold gas (T < 10^4 K) is present, so the star formation cannot proceed further. Moreover, an expanding shock front is evident (most of the gas is gathered in a narrow density strip moving outwards at high temperatures); this will result in a hot galactic wind. The hot but thin gas is being shocked during its infall, in a hot accretion mode. The low density, cold tail is formed by gas expanding and never collapsing onto the galaxy, cooling down after being photo-heated by reionizing radiation at z = 9 (see Sect. 2.3). More cold gas is present in less dense systems, both in high density regions (potentially forming new stars) and in the outskirts, the shock front being less pronounced. A large amount of dense (ρ ≈ 10^{-25} g/cm^3) and hot (T > 10^8 K) gas is also present in massive systems which are still forming stars, due to the action of feedback.

In the intermediate mass systems, a large amount of cold and dense gas is still present, ensuring a continuing star formation. In low mass systems, almost no gas is heated to T > 10^7. Thus, the gas is ejected from the galaxy at low temperatures, and no shock front develops.

Figs. 16 and 17 plot the state of the gas in two reference models, namely HDHM and LDLM; they show the radial velocity as a function of the radial distance at three different epochs and compared to the escape velocity (the blue solid lines; see the caption for more details). The mass of the ejected gas increases with time in all the models. The average velocity is proportional to the mass of the system, with massive galaxies generating winds at ≈ 2000 km/s, and low mass galaxies at about one tenth of the speed. Interestingly, the average radial velocity is proportional to the distance, like in a Hubble-like flow; the ejection of the gas is therefore an explosive phenomenon, rather than a slow outflow. However, looking more carefully it can be noticed that there are different linear velocities at the same radial distances, meaning that many explosive events have taken place. There is also a huge mass of gas moving outwards at slower speed, at late times (particularly in the HDHM model).

In Fig. 18 we plot fractional amounts of escaping gas as a function of time, for four of the models (namely, HDHM, LDHM, HDLM, LDLM; the other models show similar behaviours). We plot the trends only after the systems have virialized; in other words, the total gravitational energy must have become constant to avoid the initial phases of galaxy assembly during which the gas accumulates toward the center and the galaxy is subject to episodes of quick expansions and/or contractions. The red dotted lines display the fractional mass of gas which is found outside a given radius r_{ext} from the centers of the systems at the given time. For massive systems we adopt r_{ext} = 200 kpc, while for small systems r_{ext} = 10 kpc (these values are similar to the virial radii of the models, see Tab. 3). The blue dashed lines are relative the fraction of gas which has an outward-directed radial velocity, and is moving faster than the Hubble flow at the correspondent redshift. Finally, the black solid lines are the fraction of gas which is also above the escape velocity of the system. Clearly, two different regimes are at work. In the high mass realm, the fraction of gas which is really leaving the galaxy is only a few percent. On the contrary, in low mass systems more than half of the total gas has sufficient velocity to escape. This is easily understandable. Assuming constant density in models of different total masses, the gravitational potential energy of a system is proportional to M^{5/3}, while the energy injection from SN explosions is proportional to the number of stars in the galaxy, i.e. roughly ∝ M. Thus, the binding energy of a system grows more rapidly than the kinetic energy of its components, making it more difficult to escape from the system. However, this distinction is theoretical, and what we observe in the real Universe is simply the mass of gas that can be found outside a galaxy, moving outwards (i.e., the blue dashed line in the diagram). In this case, the situation in the two cases is much similar, however still with a higher efficiency in the low mass regime. The escaping gas is chemically enriched, as illustrated in Fig. 19 for the HDHM model.
3.5 Critical discussion of the results

Is the strong dependence of the SFH on the initial galaxy mass and density a physical result, or is it just a spurious numerical effect? The large difference in particle masses between the various simulations, a necessary choice to model galaxies spanning a wide range of sizes, is the major reason for this serious uncertainty. In high mass models, the gas particles have a mass $m_{HM} \simeq 5 \times 10^7 M_\odot$, whereas in the low mass models the mass of a gas particle is $m_{LM} \simeq 1.2 \times 10^4 M_\odot$, i.e. a factor of $\simeq 4 \times 10^3$ times smaller. Does this difference in mass resolution affect the global results of our study?

To check this issue, we run three low resolution models. The initial haloes for these test models were crudely obtained from the three high density models HDHM, HDIM, and HDLM, by simply considering one particle out four when reading the input files, and simultaneously multiplying the particle masses by a factor of four. Although this procedure is not rigorous, still it can be safely used to get a quick insight of the problem. The new models shortly indicated by lrHDHM, lrHDIM, lrHDLM have the total number of particles reduced by a factor of 4 and can be calculated on a single CPU. Their star formation is compared to that of their “parent” models and the results are plotted in Fig. 20. Even if some minor differences can be noticed, it is evident that the global trends are quite similar at the different resolutions. This is reassuring, because it implies that we can trust the results of our reference models, at least as far as their different star formation histories is concerned. This also shows that meaningful results can be obtained even using a small number of particles with much less numerical effort.

Finally, we consider the effects of imposing constant minimum threshold values to the softening and smoothing lengths of particles. To do so, we re-run the low mass models imposing minimum values $\epsilon_{\min} = h_{\min} = 100$ pc. We plot the SFHs of the resulting models in Fig. 21 and compare them with the SFHs of the same models belong to the reference set in which no limitations have been imposed to the softening and smoothing lengths. Clearly, there are non-negligible differences. The global efficiency in the star formation process is lower in the models including the threshold. Moreover and perhaps even more important, the dynamical evolution is substantially different, because in these models star formation is essentially halted after a strong peak of activity, whereas in the reference models it proceeds at nearly constant efficiency. To cast light on this issue we look at the temporal variation of the kinetic, gravitational potential, thermal, and total energies of the models under consideration. They are plotted in Fig. 22. It is clear that imposing the limit values to the softening and smoothing lengths causes an uncorrect behaviour at the epoch of the strong ejection of hot gas following the first intense burst of star formation, when the dynamical stability of the system is put in troubles. The model response is not physically adequate to drive the correct dynamical behaviour of the whole system. This means that the wrong choice of these parameters can deeply affect the overall evolution of the model. Variable softening and smoothing lengths are therefore highly recommended. Nevertheless, a great deal of the general behavior of the reference models (in particular the gross features of the star formation history) still survives also in those with the constrained softening and smoothing lengths, thus securing that even with wrong choices for these two parameters we catch the essence of the problem.
4 MASS DENSITY PROFILES

The geometrical structure of the model galaxies is best illustrated by the surface and volume density profiles. To compare these we choose a certain evolutionary stage, namely the model galaxies at a certain redshift, i.e. $z = 1$. Fig. 23 displays the surface density radial profiles projected on the $XY$ plane at such redshift. Over-plotted is the best Sersic profile,

$$
\sigma_S(r) = \sigma_0 \times (\frac{r}{r_e})^{1/m} - 1
$$

where $r_e$ is the effective radius of the galaxy, as defined e.g. in Hernquist (1990), $\sigma_0$ is the surface density at $r_e$, and $m$ is the Sersic index (if $m = 4$ the de Vaucouleurs profile is obtained). All profiles are computed starting at 0.2% of the virial radius of the galaxies to avoid the very central region of a galaxy where softening may introduce spurious numerical effects. The best-fitting Sersic index is $m \approx 4$, $m \approx 1.5$, and $m \approx 2.5$, for high-, intermediate-, and low-
mass models, respectively. In other words, high-(relative to low- and intermediate-) mass models tend to have higher \( m \), in qualitative agreement with the existence of a luminosity-Sersic index relation of ETGs (Caon et al. 1993). However, one should notice that the most massive ellipticals in the local Universe tend to have \( m \sim 8 \) (e.g. Ferrarese et al. 2006), while we find \( m \sim 4 \). Moreover, our intermediate-mass models have somewhat lower \( m \) than the low-mass ones, which is only marginally consistent with luminosity- \( m \) relation, considering its large scatter.

The overall agreement between the models and the Sersic curves is good in the external regions. However, a clear departure from the expected fits is evident in all models at small radii (some fraction of the effective radius, indicated with the black diamond shown in each panel). The central regions of the model galaxies tend to flatten out with a plateau in the mass density profile. Given the adaptiveness of the force softening, this feature can hardly be ascribed to numerical artifacts. The close scrutiny of the problem, suggests us ascribe it to the high value of the dimensionless efficiency, \( \epsilon_{\text{SF}} = 1 \), that we have adopted for the star formation rate.

This conclusion is strengthened by the three-dimensional mass density profiles shown in Fig. 24. These profiles are computed starting at 0.05% of the virial radii, and they are compared with the best fits of the Hernquist (1990) and Navarro et al. (1996) profiles (NFW), for stars and DM, respectively. While DM haloes smoothly follow the expected NFW profiles, the density of the stellar component never exceeds that of the DM, and the central regions of the model galaxies are dominated by the DM, contrary to the expectations (e.g. Tortora et al. 2009; Padmanabhan et al. 2004, but see also Grillo, 2010).

The relations between the efficiency of star formation and the gas density at which star formation is straightforward. If star formation proceeds slowly (low values of \( \epsilon_{\text{SF}} \), the gas can get high densities before it gets consumed by star formation itself. The opposite if star formation proceeds at high efficiency, high values of \( \epsilon_{\text{SF}} \). As a consequence there is an immediate correlation between star formation and the dynamical behavior of a galaxy. Allowing gas particles to turn into stars as soon as they fulfill the required numerical criteria (that is, with unit efficiency) will generally result in quite diffuse stellar systems, with shallow potential wells and large effective radii. On the contrary, a low star formation efficiency leads to dense clumps of cold gas which give birth to more compact stellar clouds (and to a delay in the beginning of the activity).

Interestingly, this puts a constraint on the modality in...
Figure 23. Stellar surface density profiles for the twelve reference models at the final time-step of each simulation. Plotted is the projected surface density $\sigma$ [g/cm$^2$] against the logarithmic radius [kpc]. Blue dots: stellar particles radial averages; dashed line: best Sersic fit (see text for details). The black diamonds are placed at the effective radii $r_e$. The red squares are placed at the smallest value reached by adaptive softening lengths, thus indicating the effective resolution of the simulations. Left to right: high density, intermediate density, low density, very low density. Top row: high mass; intermediate row: intermediate mass; bottom row: low mass.

Figure 24. Density profiles for the twelve reference models at $z \simeq 1$. Plotted is the density $\rho$ [g/cm$^3$] against the logarithmic radius [kpc]; blue dots: stellar particles; cyan dots: DM particles; dashed lines: best Hernquist and NFW fits. Left to right: high density, intermediate density, low density, very low density. Top row: high mass; intermediate row: intermediate mass; bottom row: low mass.

which gas collapses and forms stellar systems. Historically, there are two possible and competing theoretical scenarios to describe the assembly of the stellar mass of an elliptical galaxy (excluding the possibility of merger events). In the first one, stars originally form far away from the effective radius of the galaxy and subsequently accrete onto its inner regions, in a non-dissipative fashion. In the other case, first the gas flows into the central regions and then it is turned into stars. If the correct scenario is the first one, smoothing of the central density cusp of the profiles is expected, because the infalling stellar clumps loose orbital energy through dynamical friction and the halo heats up [Bertin et al. 2003]. In the second case, the central distribution is expected to steepen up, because the gas radiates away its internal energy and an adiabatic contraction takes place [Gnedin et al. 2004]. Obviously, both processes are expected to play a role...
in reality, and this is indeed what we see in our models, in which however the final density profiles show a plateau in the central regions, as expected in the “infalling stars” model.

Our results also clearly suggest that if stars are formed too early, the resulting stellar systems are too shallow, ending up in “wrong” profiles: stars will tend to conserve their velocity dispersion as they fall into the potential wells and will therefore have elongated orbits, which is not the case when gas is let dissipate part of its kinetic energy and reach very high densities before turning into stars.

In this context, it is worth recalling that there is some observational evidence for a deficit in the central luminosity of a few early type galaxies as recently found in observational data by Côté et al. (2007, their Fig. 1), and references therein). However, the observed flattening begins at smaller radii (a few percent of $r_c$), and it is present only in bright (i.e. massive) systems, whereas the small systems seem to present a luminosity excess. Indeed, looking at the panels in the bottom row of Fig. 23 one can notice that the central profiles of low mass models show some scatter and tend to steepen in the very central regions (especially in the VLDLM model). Recalling that the observational information stands on the luminosity profile, given the presence of young and hence luminous stellar populations in the cores of our low mass galaxy models, (and their absence in the cores of the massive ones), we can speculate that the trends of the observed curves should be fairly reproduced by translating the mass profiles into luminosity into luminosity profiles. This is an interesting point, because the excess or deficit of light could be explained just in terms of the stars and DM orbits, without the presence of a super-massive black hole affecting the global dynamics in the center of the systems, as instead suggested in the studies quoted above.

5 THE MASS-RADIUS RELATION

To conclude our analysis, we briefly consider the Mass-Radius Relation (MRR) for ETGs. An overview of recent results concerning estimates of the star mass and dimensions of ETGs is given by Burstein et al. (1997) which displays the sample of ETGs collected by Bernardi et al. (2010, green dots). The sample contains about 65,000 objects. We have also added the small group of galaxies (about 600 objects in total) studied long ago by Bernardi et al. (2010). The green solid line is their linear best fit, log $r_{1/2} = 0.54 \log M_s - 5.25$. The contour lines display the regions of the MR-plane populated by the same number of objects (the bar yields the correspondence between color-code and number of galaxies). The data are from the HB sample of Bernardi et al. (2010). The green solid line is their linear best fit, log $r_{1/2} = 0.54 \log M_s - 5.25$. The black diamonds are the data from Burstein et al. (1997), which also include the dwarf galaxies. Large squares: the twelve major models at the last simulated age; the same colors refer to same initial densities. Black: high density. Blue: intermediate density. Magenta: low density. Red: very low density. The thin dashed black lines are fits to the models (see text).

![Figure 25. Mass-radius relation, i.e. the relation $\log(r_{1/2})$ versus $\log M_s$, where $r_{1/2}$ is the half mass (light) radius, and $\log M_s$ is the total stellar mass in units of $10^{12} M_\odot$. The contour lines display the regions of the MR-plane populated by the same number of objects (the bar yields the correspondence between color-code and number of galaxies). The data are from the HB sample of Bernardi et al. (2010). The green solid line is their linear best fit, log $r_{1/2} = 0.54 \log M_s - 5.25$. The black diamonds are the data from Burstein et al. (1997), which also include the dwarf galaxies. Large squares: the twelve major models at the last simulated age; the same colors refer to same initial densities. Black: high density. Blue: intermediate density. Magenta: low density. Red: very low density. The thin dashed black lines are fits to the models (see text).](image-url)

for. However, the thorough analysis of this issue made by Chiosi et al. (2012a) has shown that the observational MRR is the result of a subtle interaction among different concurring physical causes.

We begin by noting that as expected models with equal initial over-densities fall along the straight lines $\log r_{1/2} = \alpha \log M_s + \beta$, with slope $\alpha$ ranging from 0.2 to 0.3 and $\beta$ depending on the redshift $\beta \propto (1 + z)^{-1}$. Slopes close to 0.33 are indeed what is predicted for collapsing spherical perturbations of DM (see e.g. Fan et al. 2010). The deviation from the pure $\log r_{1/2} = \alpha \log M_s + \beta$ relationship is larger towards galaxies of lower mass and it is due to the complex baryon physics; the discussion by Chiosi & Carraro (2002) and Chiosi et al. (2012a) on the issue still applies also here. Chiosi et al. (2012a) argue that the observed MRR for ETGs is the result of two complementary mechanisms. On one hand, the mass function of DM haloes hosting the visible galaxies gives (i) the typical cut-off mass at which, at any redshift, haloes become “common” on a chosen spatial scale, and (ii) the typical epoch at which low mass haloes begin to vanish at a rate higher than that at which they are born because of merger events (Lukić et al. 2007). On the other hand, these constraints define two loci (curves) on the MR-plane, because to each mass and redshift a typical dimension (i.e., radius) can be associated (using the basic relation between mass and radius of a collapsing object).
If the typical dimension of a galaxy is somehow related to that of the hosting DM halo (as our NB-TSPH models seem to suggest), then the region of the MR-plane between the two limits fixed by the halo mass function is populated by galaxies whose dimensions are fixed at the epoch of formation, and only those objects that are “possible” at any given epoch may exist, populating a narrow region of the MR-plane. All details of this rather complex analysis can be found in Chiosi et al. (2012a) to whom the reader should refer.

## 6 SUMMARY AND CONCLUSIONS

We have presented a numerical investigation on the star formation histories of ETGs, performed with the aid of hydrodynamical NB-TSPH simulations of galaxy formation in cosmological context. By varying the initial properties of a single model in terms of total mass and initial over-density, we investigated how the initial configuration influences the formation of the system along the Hubble time.

The main result of our study is that the star formation history of a galaxy is directly dependent on its initial total mass and initial over-density. Massive galaxies ($M_{\text{tot}} \approx 10^{13} M_\odot$) build their stellar content passing through an initial burst of activity at early times. In contrast, low mass systems ($M_{\text{tot}} \approx 10^9 M_\odot$) continue to form stars throughout their lifetime, with episodes of intense activity separated by period of relative quiescence. Moreover, galaxies with the same initial mass have different star formation histories depending on their initial over-density: the deeper the perturbation, the more peaked, early and intense is the star forming activity. This is particularly true for intermediate mass galaxies ($M_{\text{tot}} \approx 10^{11} M_\odot$). Thus, denser objects, which may be identified with cluster and group central galaxies, have more peaked and earlier burst of star formation, while less dense objects (i.e., field and satellite galaxies) usually have delayed and lower activity. This is largely consistent with previous analyses by Chiosi & Carraro (2002), Clemens et al. (2006), and Thomas et al. (2005) to mention a few.

The structural properties of our models are in good agreement with those of the observed real galaxies, with some important differences (e.g. the density profiles in the central regions, the mass-to-metallicity relation) which we ascribe to the unusual choice we have made for the star formation dimensionless efficiency parameter, $\epsilon_{\text{SF}} = 1$.

The picture emerging from this study is that an early-hierarchical, monolithic-like intense activity of star formation at high redshift is possible starting from physically sounded cosmological conditions, and is fully adequate to explain the observed features of massive elliptical galaxies.

Major mergers are certainly possible, but they are not strictly required. It is indeed easy to understand that on the contrary they would alter the subtle game among different physical ingredients all contributing to shape the body of properties and their mutual tight correlations shown by early type galaxies such as for instance the MRR itself (Chiosi & Carraro 2002, Chiosi et al. 2012). Along this line of thought Buonomo et al. (2000) showed how a major merger would shift the position of the parent objects towards the upper region of the MR plane, where no observed galaxy is found. Moreover, chemical features of the elliptical galaxies cannot be reproduced if (wet) merging events are considered. Finally, the shape of the galaxy mass function Bundy et al. (2006, Pérez-González et al. 2008) rules out the possibility that late major dry mergers are responsible for the mass assembly of massive systems, out of smaller but already passively evolving objects. Our models show instead that massive galaxies can form out of single, peaked perturbations, consistently with the cosmological background.

Another interesting point to note is that the stellar feedback seems to be sufficient to quench the star formation process in massive objects without invoking more exotic sources of energy, like Active Nuclei, provided it is modeled with sufficient accuracy. Since the models are evolved in void, the lack of late-time infall of gas could explain why the presence of an AGN is not strictly necessary. In relation to this, one could argue that strong galactic winds moving outward at high velocity are probably sufficient to inhibit flows of material towards the inner regions. However, the occurrence of Active Nuclei cannot be excluded.

Furthermore, we are also able to recover the complex, episodic star formation histories typical of many dwarf galaxies. The galactic breathing phenomenon is a natural consequence of an accurate treatment of the stellar feedback process.

We conclude that the fate of a proto-galaxy is essentially determined by its initial conditions in terms of mass and over-density, while external factors such as encounters, mergers and disruptions, while substantially altering the evolution of the involved systems, are not a fundamental ingredient in the global evolution of the early type galaxies population.

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