Zooming on the internal structure of $z \simeq 6$ galaxies

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ABSTRACT
We present zoom-in, AMR, high-resolution ($\simeq 30$ pc) simulations of high-redshift ($z \simeq 6$) galaxies with the aim of characterizing their internal properties and interstellar medium. Among other features, we adopt a star formation model based on a physically sound molecular hydrogen prescription, and introduce a novel scheme for supernova feedback, stellar winds and dust-mediated radiation pressure. In the zoom-in simulation the target halo hosts “Dahlia”, a galaxy with a stellar mass $M_* = 1.6 \times 10^{10} M_\odot$, representative of a typical $z \sim 6$ Lyman Break Galaxy. Dahlia has a total $H_2$ mass of $10^{8.5} M_\odot$, that is mainly concentrated in a disk-like structure of effective radius $\simeq 0.6$ kpc and scale height $\simeq 200$ pc. Frequent mergers drive fresh gas towards the centre of the disk, sustaining a star formation rate per unit area of $\simeq 15 M_\odot$ yr$^{-1}$ kpc$^{-2}$. The disk is composed by dense ($n \gtrsim 25$ cm$^{-3}$), metal-rich ($Z \simeq 0.5 Z_\odot$) gas, that is pressure-supported by radiation. We compute the 158 nm $[C\,\text{II}]$ emission arising from Dahlia, and find that $\simeq 95\%$ of the total $[C\,\text{II}]$ luminosity ($L_{[C\,\text{II}]} \simeq 10^{5.5} L_\odot$) arises from the $H_2$ disk. Although 30\% of the $C\,\text{II}$ mass is transported out of the disk by outflows, such gas negligibly contributes to $[C\,\text{II}]$ emission, due to its low density ($n \lesssim 10$ cm$^{-3}$) and metallicity ($Z \lesssim 10^{-1} Z_\odot$). Dahlia is under-luminous with respect to the local $[C\,\text{II}]-SFR$ relation; however, its luminosity is consistent with upper limits derived for most $z \sim 6$ galaxies.

Key words: galaxies: high-redshift, formation, evolution, ISM – infrared: general – methods: numerical

1 INTRODUCTION
The discovery and characterization of primeval galaxies constitute some of the biggest challenges in current observational and theoretical cosmology.

Deep optical/near infrared (IR) surveys (Dunlop 2013; Madau & Dickinson 2014; Bouwens et al. 2015) have made impressive progresses in identifying galaxies well within the Epoch of Reionization ($z \simeq 6$). Such surveys yield key information about the star formation (SF) of hundreds of galaxies in the early Universe. They also allow to statistically characterize galaxies in terms of their UltraViolet (UV) luminosity up to $z \sim 10$ (Bouwens et al. 2015). However – using these surveys broad band alone – little can be learned about other properties as their gas and dust content, metallicity, interactions with the surrounding environment (e.g. Barnes et al. 2014), feedback (e.g. Dayal et al. 2014), and outflows (Gallerani et al. 2016).

To obtain a full picture of these systems, optical/IR surveys must be complemented with additional probes. Information on the metal content and energetics of the interstellar medium (ISM) can be obtained with observations of Far IR (FIR) fine structure lines, and in particular the $[C\,\text{II}] (^2P_{3/2} \rightarrow ^2P_{1/2})$ line at 157.74 µm. The $[C\,\text{II}]$ line is the dominant coolant of the ISM being excited in different ISM phases, as the diffuse cold neutral medium (CNM), warm neutral medium (WNM), high density photodissociation regions (PDRs), and – to a lower extent – ionized gas (Tielens & Hollenbach 1985; Wolfire et al. 1995; Abel 2006; Vallini et al. 2013). As $[C\,\text{II}]$ emission can be enhanced by

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shocks, it has been suggested as a good outflow tracer (e.g. Maiolino et al. 2012; Kreckel et al. 2014; Cicone et al. 2015; Janssen et al. 2016), and can thus in general be used to study feedback processes in galaxies.

Observationally, the [C\textsc{ii}] line is a promising probe as it is often the brightest among FIR emission lines, accounting for up to \(\sim 1\%\) of the total IR luminosity of galaxies (e.g. Crawford et al. 1985; Madden et al. 1997). It has been successfully used to probe the low-z ISM (e.g. De Looze et al. 2014). The unprecedented sensitivity of the Atacama Large Millimeter/Submillimeter Array (ALMA) makes it possible for the first time to use [C\textsc{ii}] emission to characterize high-z galaxies. Before the ALMA advent, in fact, detections were limited to a handful of QSO host galaxies, and rare galaxies with extreme SF rates (\(SFR \simeq 10^6 M_\odot \text{yr}^{-1}\)) at \(z \sim 6\) – 7 early ALMA searches for [C\textsc{ii}] lines have mostly yielded upper limits (e.g. Ouchi et al. 2013; Kanekar et al. 2013; Ota et al. 2014; Schaefer et al. 2015). The situation has changed recently with a number of robust [C\textsc{ii}] detections (e.g. Maiolino et al. 2015; Capak et al. 2015; Willott et al. 2015; Knudsen et al. 2016).

However, for “normal” star forming galaxies (\(\lesssim 10^6 M_\odot \text{yr}^{-1}\)) at \(z \sim 6\) – 7, early ALMA searches for [C\textsc{ii}] lines have mostly yielded upper limits (e.g. Ouchi et al. 2013; Kanekar et al. 2013; Ota et al. 2014; Schaefer et al. 2015). The situation has changed recently with a number of robust [C\textsc{ii}] detections (e.g. Maiolino et al. 2015; Capak et al. 2015; Willott et al. 2015; Knudsen et al. 2016).

In many cases the high-z [C\textsc{ii}] line luminosity is fainter than expected from the [C\textsc{ii}]-SFR relation found in local galaxies (De Looze et al. 2014). To explain such [C\textsc{ii}]-SFR deficit, some efforts have been devoted to model the [C\textsc{ii}] emission from high-z galaxies (Nagamine et al. 2006; Vallini et al. 2013; Muñoz & Furlanetto 2014; Vallini et al. 2015; Olsen et al. 2015). In brief, these theoretical works show that the [C\textsc{ii}]-SFR deficit can be ascribed to different effects:

\begin{itemize}
  \item[(a)] Lower metallicity of high-z galaxies (Vallini et al. 2013; Muñoz & Furlanetto 2014; Vallini et al. 2015), in particular supported by observations of lensed galaxies (Knudsen et al. 2016).
  \item[(b)] Suppression of [C\textsc{ii}] line around star forming regions Vallini et al. (2013), typically observed as a displacement of the [C\textsc{ii}] with respect to the UV emitting region, as seen e.g. in BDF3299 (Maiolino et al. 2015) and in some of the Capak et al. (2015) galaxies. This would be a signature of the collisional ionization of the [C\textsc{ii}]-emitting gas.
  \item[(c)] Suppression of [C\textsc{ii}] line by the increased CMB temperature in the WNM/CNM component (Pallottini et al. 2015; Vallini et al. 2015), similarly to what observed for dust emission (da Cunha et al. 2013).
\end{itemize}

Simulating the ISM of early galaxies at sufficient resolution and including feedback effects might shed light on these questions. Feedback prescriptions are particularly important as such process regulates the amount of (dense) gas likely to radiate most of the power detected with FIR lines. Several studies have explored optimal strategies to include feedback in galaxy simulations.

For some works, the interest is in the comparison between different kind of stellar feedback prescription, as modelled via thermal and/or kinetic energy deposition in the gas from supernovae (SN), winds (Agertz et al. 2013; Hopkins et al. 2014; Barai et al. 2015; Agertz & Kravtsov 2015), and thus to the Poisson equation on the AMR

\begin{table}
\centering
\begin{tabular}{|c|c|c|c|c|}
\hline
\text{cosmo} & \text{zoom} & \text{\(\Delta_m^{\text{max}}\)} & \text{\(\Delta_m^{\text{min}}\)} & \text{\(\Delta_x^{\text{min}}\) at \(z = 6\)} \\
\hline
\text{3.4 \times 10^7} & \text{6.7 \times 10^4} & \text{78.1} & \text{78.1} & \text{2.5 \times 10^3} \\
\hline
\text{78.1} & \text{1.2 \times 10^4} & \text{9.7} & \text{0.1} & \text{32.1} \\
\hline
\end{tabular}
\caption{Resolution setup for the cosmological run (cosmo) and subsequent zoom-in (zoom) simulation. \(m_{dm}\) and \(m_\delta\) are in units of \(M_\odot/h\) and indicate the dark matter (DM) and baryon mass resolution, respectively; \(\Delta_M^{\text{max}}\) and \(\Delta_M^{\text{min}}\) indicate the coarse grid and minimum available refinement scale, respectively. Both scales are reported in comoving kpc/h. For \(\Delta_M^{\text{min}}\) we also report the maximum physical pc scale at \(z = 6\). For the cosmo run, no refinement is used, and for the zoom, we indicate the increased resolution of the zoomed halo due to the multi-mass approach and the AMR.}
\end{table}

other analyses focus on implementing complex chemical networks in simulations (Tomassetti et al. 2015; Maio & Tescari 2015; Bovino et al. 2016; Richings & Schaye 2016; Grassi et al. 2016), radiative transfer effect (Petkova & Maio 2012; Rosdahl et al. 2015; Maio et al. 2016), or aim at removing tensions between different coding approaches (Kim et al. 2014).

Thus, we can improve galaxy simulations by providing theoretical expectations for [C\textsc{ii}] that should be compared with state-of-the-art data. Such a synergy between theory and observations, in turn, can guide the interpretation of upcoming ALMA data and drive future experiments of large scale [C\textsc{ii}] mapping (Gong et al. 2012; Silva et al. 2015; Yue et al. 2015; Pallottini et al. 2015), which would lead to a characteristic high-z galaxy population. In the present work we simulate a galaxy detected in [C\textsc{ii}] with ALMA current observations.

The paper is structured as follows. In Sec. 2 we detail the numerical model used to set-up the zoom-in simulation, and describe the adopted H\textsubscript{2} star formation prescription (Sec. 2.2), mass and energy inputs from the stellar populations (Sec. 2.3) and feedback (including SN, winds and radiation pressure) (Sec. 2.4). The results are discussed in Sec. 3, where we analyze star formation history and feedback effects in relation to ISM thermodynamics (Sec. 3.2) and its structural properties. The expected [C\textsc{ii}] emission and other observational properties of high-z galaxies are discussed in Sec. 3.3. Conclusions are given in Sec. 4.

2 NUMERICAL SIMULATIONS

We carry out our simulation using a customized version of the adaptive mesh refinement (AMR) code RAMSES (Teyssier 2002). RAMSES is an octree-based code that uses Particle Mesh N-body solver for the dark matter (DM) and an unsplit 2nd-order MUSCL\textsuperscript{2} scheme for the baryons. Gravity is accounted for solving the Poisson equation on the AMR grid via a multi-grid scheme with Dirichlet boundary conditions on arbitrary domains (Guillemet & Teyssier 2011). For the present simulation we choose a refinement based on a Lagrangian mass threshold-based criterion.

\begin{itemize}
  \item[(2)] MUSCL: Monotone Upstream-centred Scheme for Conservation Laws
\end{itemize}
Chemistry and heating/cooling processes of the baryons are implemented with GRACKELE.4 (Bryan et al. 2014), the standard library of thermo-chemical processes of the AGORA project (Kim et al. 2014). Via GRACKELE, we follow the H and He primordial network and tabulated metal cooling and photo-heating rates calculated with CLOUDY (Ferland et al. 2013). Cooling includes also inverse Compton off the cosmic microwave background (CMB), and heating from a redshift-dependent ionizing UV background (Haardt & Madau 2012, UVB). Since H2 gas phase formation is not accounted for, we do not include the cooling contribution of such species.

Because of stellar feedback (Sec 2.3 and 2.4), the gas can acquire energy both in thermal and kinetic form. The distinction is considered by following the gas evolution of the standard thermal energy and a “non-thermal” energy (Agertz & Madau 2012, UVB). Since H2 gas phase formation is not accounted for, we do not include the cooling contribution of such species.

2.1 Initial conditions

The initial conditions (IC) used for the suite are generated with MUSIC (Hahn & Abel 2011). MUSIC produces IC on nested grid using a real-space convolution approach (cf. Bertschinger 1995). The adopted Lagrangian perturbation theory scheme is perfectly suited to produce IC for multi-mass simulations and – in particular – zoom simulations. To generate the ICs, the transfer functions are taken from (Eisenstein & Hu 1998).

To set-up the zoom-in simulation, we start by carrying out a cosmological DM-only run. The simulation evolves a volume \( V_{\text{cosmo}} = (20 \text{ Mpc}/h)^3 \) from \( z = 100 \) to \( z = 6 \) with DM mass resolution of \( m_{\text{dm}}^{\text{cosmo}} = 3.4 \times 10^7 h \text{M}_\odot \). The resolution of the coarse grid is \( \Delta x_{\text{cosmo}} = 78.1 h \text{kpc} \), and we do not include additional levels of refinement. Using HOP (Eisenstein & Hut 1998) we find the DM halo catalogue at \( z = 6 \). The cumulative halo mass function extracted from the catalogue is in agreement with analytical expectations (e.g. Sheth & Tormen 1999), within the precision of halo-finder codes (e.g. Knebe et al. 2013).

From the catalogue we select a halo with DM mass \( M_h \approx 10^{11}/h \text{M}_\odot \) (resolved by \( 5 \times 10^4 \) DM particles), whose virial radius is \( r_{\text{vir}} \approx 15 \text{kpc} \) at \( z = 6 \). Using hop we select the minimum ellipsoid enveloping \( 10 r_{\text{vir}} \), and trace it back to \( z = 100 \). As noted in Oñorbe et al. (2014), this is usually sufficient to avoid contamination. At \( z = 100 \) the trace back volume is \( V_{\text{zoom}} \approx (2.1 \text{ Mpc}/h)^3 \). Using MUSIC we recalculate the ICs, by generating 3 additional level of refinement. For such multi-mass set-up, the finer DM resolution is \( m_{\text{dm}}^{\text{zoom}} = 6.7 \times 10^3/h \text{M}_\odot \), that corresponds to a spatial resolution of \( \Delta x_{\text{zoom}} = 9.7 h \text{kpc} \). We note that because of the traced back volume, our simulation is expected to probe not only the target halo, but also its satellites and environment, similar to other works (e.g. Fiocconi et al. 2015, where the target halo is chosen at \( z \approx 3 \)).

In the zoom-in simulation \( \Delta x_{\text{zoom}} \) corresponds to our coarse grid resolution, and we allow for 6 additional refinement levels, based on a Lagrangian mass threshold-based criterion. At \( z = 6 \), the baryonic component of the selected halo has a mass resolution of \( m_b = 1.8 \times 10^4 h \text{M}_\odot \) and a physical resolution of \( \Delta x_{\text{min}} = 31.9 h \text{pc} \). For convenience, a summary of the resolution outline can be found in Tab. 1. Note that the refined cell of our simulations have mass and size typical of molecular clouds (MC, e.g. Gorti & Hollenbach 2002; Federrath & Klessen 2013).

In the present paper we refer to metallicity (Z) as the sum of all the heavy element species without differentiating among them, and assume solar abundance ratios (Asplund et al. 2009). In the IC, the gas is characterized by a mean molecular weight \( \mu = 0.59 \), and has metallicity \( Z = Z_{\text{floor}} > 0 \). The metallicity floor mimics the pre-enrichment of the halo at high-z, when we do not have the resolution to follow precisely star formation and gas enrichment. We set \( Z_{\text{floor}} = 10^{-2} Z_{\odot} \), a level that is compatible with the metallicity found at high-z in cosmological simulations for diffuse enriched gas (Davé et al. 2011; Palla et al. 2014; Maio & Tescari 2015). Note that such low metallicity only marginally affects the gas cooling time, but is above the critical metallicity for formation of Population III stars. Additionally, a posteriori, we have found that the metallicity floor contribute for only \( \lesssim 0.2\% \) of the total metal mass produced by stars by \( z = 6 \) in the refined region.

2.2 Star formation model

We model star formation (SF) by assuming a H2 dependent Schmidt-Kennicutt relation (Schmidt 1959; Kennicutt 1998)

\[
\dot{\rho}_* = f_{\text{H}_2} \rho / t_{\text{sf}}
\]

where \( \dot{\rho}_* \) is the local SF rate (SFR) density, \( f_{\text{H}_2} \) the molecular hydrogen fraction, \( \rho \) the gas density and \( t_{\text{sf}} \) the SF time scale. In eq. 1a we assume the SF time scale to be proportional to the free-fall time, i.e.

\[
t_{\text{sf}} = c_{\text{sf}}^{-1} \sqrt{3\pi/(32 G \rho)} ,
\]

where \( c_{\text{sf}} \) describes the SF efficiency and it is treated as a parameter in the present work (cf. Semenov et al. 2016, see dis-

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3 See also https://grackle.readthedocs.org/

4 While the distinction in thermal and non-thermal is similar to previous works (e.g. Agertz & Kravtsov 2015), we note that usually the time scale for dissipation is fixed to 10 Myr.

5 A posteriori, we have checked that the halos in the zoom-in region have a contamination level \( \lesssim 0.1\% \).
H₂ fraction \( f_{H_2} \) as a function of gas density \( n \) obtained using the KTM09 model (eqs. 2). Different solid lines correspond to different metallicity \( Z \) of the gas. Horizontal dotted grey lines mark \( f_{H_2} \) values of 0.5 and 1. Vertical dashed lines indicate the critical density \( n_c \) where \( f_{H_2} = 0.5 \) for different \( Z \); these critical density values are obtained as a fit (eq. 3b) to the KTM09 model (eq. 2). See the text for details.

As shown in (Krumholz & Gnedin 2011), for \( Z \gtrsim 10^{-2}Z_\odot \) such approximation gives H₂ fractions compatible with those resulting from a full non-equilibrium radiative transfer calculations.

In Fig. 1 we plot \( f_{H_2} \) from the KTM09 model as a function of the gas density. Different solid lines refer to different metallicity. At a fixed metallicity, the molecular fraction as a function of density vanishes for low values of \( n \); it steeply rises up to \( f_{H_2} \sim 0.8 \) in one density dex and asymptotically reaches \( f_{H_2} = 1 \). The critical density where the gas can be considered molecular \( (f_{H_2} = 0.5) \) is roughly inversely proportional to the metallicity, i.e. \( n_c \sim 25(Z/Z_\odot)^{-1} \text{cm}^{-3} \) (see also Agertz et al. 2013). We note that when detailed chemistry calculations are performed, such critical density depends on the chemical network and the assumptions regarding gas shielding from external radiation and clumpiness. As a consequence, the actual critical density can be higher that the one predicted by the KTM09 model (e.g. Bovino et al. 2016).

Because of the particular shape of the \( f_{H_2}(n) \) relation, the adopted SF law (eqs. 1–2) is roughly equivalent to a prescription based on a density threshold criterion:

\[
\rho_* = \Theta(n - n_c) m_p n / \bar{m}_p ,
\]

where \( m_p \) is the proton mass and the critical density

\[
\n_c \simeq 26.45 \left( \frac{Z}{Z_\odot} \right)^{-0.87} \text{cm}^{-3}
\]

is calculated as a fit to the \( f_{H_2} \), KTM09 model. In Fig. 1, we show \( n_c \) for various metallicities (dashed vertical lines).

Eqs. 3 are not used to calculate the \( SFR \) in the simulation. However, being simpler, such formulation can be used to enhance our physical intuition of the adopted SF law\(^6\) in analyzing the results. As noted in Hopkins et al. (2013), the morphology of a galaxy is very sensitive to the minimum density of the cells that are able to form star.

During the simulation, eqs. 1 are solved stochastically, by drawing the mass of the new star particles from a Poisson distribution (Rasera & Teyssier 2006; Dubois & Teyssier 2008; Pallottini et al. 2014a). We impose that no more than half of a cell mass can be converted into a star particle in each event. This prescription ensures the numerical stability of the code (Dubois & Teyssier 2008). This is also consistent with the picture that nearly half of the mass in a MC is Jeans unstable (Federrath & Klessen 2013).

We allow SF only if the mass of a new star particle is at least equal to the baryon mass resolution. This avoids numerical errors for the star particle dynamics and enables us to treat the particle as a stellar population with a well sampled initial mass function (IMF). Additionally, the SF law is driven by H₂ formation on dust grains, we do not allow gas to form stars if the dust temperature is larger than \( \simeq 2 \times 10^4 \), because of dust sublimation (see Sec. 2.4 and App. B for the details on the dust prescriptions).

For the present work we assume a SF efficiency \( \eta_{sf} = 10\% \), in accordance with the average values inferred from

\(6\) As a consequence of the rough equivalence, it is not necessary to manually prevent SF in underdense regions, by imposing that an overdensity \( \Delta > 200 \) is needed to form stars. At the start of the simulation \( (z = 100) \), the mean density of the gas is \( \sim 0.1 \text{mp}_c \text{cm}^{-3} \), while the “effective” SF threshold would be \( n_c \sim 10^4 \text{cm}^{-3} \) for gas at \( Z = Z_{\text{floor}} \).
Because of the finite mass resolution, it is necessary to introduce (according to eqs. 1–2d) “star particles” to represent stellar populations. To this aim, we adopt a Kroupa (2001) IMF

\[ \Phi(m) \propto \left[ m^{-\alpha_1} \Theta(m_1 - m) + m^{-\alpha_2} \Theta(m - m_1) m_1^{\alpha_2 - \alpha_1} \right], \]

where \( \alpha_1 = 1.3, \alpha_2 = 2.3, m_1 = 0.5 M_\odot \), and \( m \) is in the range \([10^{-1} - 10^{3}] M_\odot \). The proportionality constant is chosen such that

\[ \int_{0.1 M_\odot}^{100 M_\odot} m \Phi \, dm = 1. \]

Once formed, stars affect the environment with chemical, mechanical and radiative feedback. These stellar inputs are parameterized by the cumulative fraction of the returned gas mass, metals and energy (e.g. Salvadori et al. 2008; de Bennasuti et al. 2014; Salvadori et al. 2015). Mass and energy inputs are conveniently expressed per unit stellar mass formed (\( M_* \)).

Chemical feedback depends on the return fraction (\( R \)) and the yield (\( Y \)):

\[ R(t_s) = \int_{m(t_s)}^{100 M_\odot} (m - w) \Phi \, dm \quad (5a) \]

\[ Y(t_s) = \int_{m(t_s)}^{100 M_\odot} m_z \Phi \, dm, \quad (5b) \]

where \( w(m, Z) \) and \( m_z(m, Z) \) are the stellar remnant and the metal mass produced for a star of mass \( m \) and metallicity \( Z \) (e.g. Woosley & Weaver 1995; van den Hoek & Groenewegen 1997), and \( m(t_s) \) is the minimum stellar mass with lifetime \( \tau_* \) shorter than \( t_s \), the time elapsed from the creation of the stellar particle (i.e. the “burst age”).

This approach is used both in zoom galaxy simulations (e.g. Kim et al. 2014) and cosmological simulations (e.g. P results in the ISM).

The mechanical energy input includes SN explosions and winds, either by OB or AGB stars in young (\(< 40 \) Myr) or evolved stellar populations:

\[ \epsilon_{sn}(t_s) = \int_{m(t_s) > 8 M_\odot}^{40 M_\odot} \epsilon_{snw} \Phi \, dm, \quad (5c) \]

\[ \epsilon_{w}(t_s) = \int_{m(t_s)}^{100 M_\odot} \epsilon_{w} \Phi \, dm, \quad (5d) \]

where \( \epsilon_{sn} = \epsilon_{sn}(m, Z) \) and \( \epsilon_{w} = \epsilon_{w}(m, Z) \) are the energy released by SN and stellar winds in units of \( 10^{51} \) erg. We have further assumed that only stars with \( 8 \leq m/M_\odot \leq 40 \) can explode as SN.

Radiative energy inputs can be treated within a similar formalism. The cumulative energy \( \epsilon_{12} \) associated to the spectral range \( (\lambda_1, \lambda_2) \) can be written as

\[ \epsilon_{12}(t_s) = \int_{0}^{t_s} \int_{m(t)}^{100 M_\odot} L_{12} \Phi \, dm \, dt \quad (5e) \]

\[ L_{12}(t) = \int_{\lambda_1}^{\lambda_2} L_{\lambda} \, d\lambda, \quad (5f) \]

where \( L_{\lambda} = L_{\lambda}(m, Z_s) \) is the luminosity per unit wavelength and mass. For convenience, we express the radiation energy in units of foe, as for the mechanical energy (eqs. 5c). In the following we specify \( \epsilon_{12} \) in eq. 5e, by separately considering ionizing radiation \( (\lambda_1 = 0, \lambda_2 = 912 \) Å) denoted by \( \epsilon_{ion} \), and the soft UV band, \( \epsilon_{uv} \), defined as the range \( (\lambda_1 = 912 \) Å, \( \lambda_2 = 4000 \) Å).

In eqs. 5, the quantities \( w, m_z, \epsilon_{sn}, \epsilon_w, \) and \( L_{\lambda} \) can be calculated from stellar evolutionary models. We adopt the padova (Bertelli et al. 1994) stellar tracks for metallicities \( Z_s/Z_\odot = 0.02, 0.2, 0.4, \) and 1 to compute the chemical, mechanical and radiative inputs using starburst99 (Leitherer et al. 1999, 2010).

In Fig. 2 we plot \( R, Y, \epsilon_{sn}, \epsilon_w, \epsilon_{ion} \) and \( \epsilon_{uv} \) as a function of \( t_* \). For each curve the shaded regions indicate the \( 0.02 \leq Z_s/Z_\odot \leq 1 \) metallicity range; single \( Z_s \) tracks are indicated with dark lines. The time interval during which massive stars can explode as SN (0.8 \( \gtrsim \) log \( t_*/\)Myr \( \gtrsim \) 1.6) is highlighted with vertical dashed lines, and the upper axis is labelled with the corresponding stellar mass.

Note that the OB stars contribution (log \( t_*/\)Myr \( \gtrsim \) 0.8) to \( \epsilon_{sn} \), \( Y \) and \( R \) is roughly proportional to \( t_* \) and \( Z_s \) (see also Agertz et al. 2013, in particular eqs. 4). As in the simulation the metallicity floor is set to \( Z_{floor} = 10^{-3} Z_\odot \), we slightly overestimate the wind contribution for low \( Z_s \).

Finally, note that the change of behavior of \( \epsilon_w \) at log \( t_*/\)Myr \( \lesssim \) 2 is due to the ionizing \( (\lambda \lesssim 912 \) Å) photon production suppression. At late times (log \( t_*/\)Myr \( \gtrsim \) 1.6), AGB stars give a negligible mechanical energy contribution \( (\epsilon_w \approx \) constant) but return mass and metals to the gas (\( R, Y \)).

### 2.4 Stellar feedback

Eqs. 5 provide us with the energy produced by stars in different forms. The next step is to understand what fraction of that energy is eventually deposited in the ISM. Consider a stellar population of initial mass \( M_* \), metallicity \( Z_\odot \), and age \( t_* \) residing in a gas cell with volume \( V_{cell} \). In our scheme,

\[ \epsilon_{sn} = \epsilon_{sn}(m, Z) \]

\[ \epsilon_{w} = \epsilon_{w}(m, Z) \]

where \( \epsilon_{sn} = \epsilon_{sn}(m, Z) \) and \( \epsilon_{w} = \epsilon_{w}(m, Z) \) are the energy released by SN and stellar winds in units of \( 10^{51} \) erg. We have further assumed that only stars with \( 8 \leq m/M_\odot \leq 40 \) can explode as SN.

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7 Stellar lifetimes are roughly independent of metallicity for \( Z_s > 10^{-3} Z_\odot \) (Raiteri et al. 1996, see eq. 3).

8 Similarly to Kim et al. (2014), when computing the yields in eq. 5a, we assume that the metal mass is linked to the oxygen and iron masses via \( m_Z = 2.09 m_O + 1.06 m_F \), as appropriate for Asplund et al. (2009) abundances.
when the simulation evolves for a time $\Delta t$, the chemical feedback act as follows:

$$
\rho = \rho + [R(t_s + \Delta t) - R(t_s)] M_*/V_{cell} \quad (6a)
$$

$$
Z = Z + [Y(t_s + \Delta t) - Y(t_s)] M_*/V_{cell} \quad (6b)
$$

where $\rho$ and $Z$ are the gas density and metallicity and $R$ and $Y$ are taken from eqs. 5a. Note that chemical enrichment is due both to the SN and AGB winds.

### 2.4.1 Supernova explosions

For the mechanical feedback, let us first consider the case of SNe. At each SN event the specific energy of the gas changes as

$$
e_{th} = e_{th} + f_{th} [e_{sn}(t_s + \Delta t) - e_{sn}(t_s)] M_*/V_{cell} \quad (6c)
$$

$$
e_{nth} = e_{nth} + f_{kn} [e_{sn}(t_s + \Delta t) - e_{sn}(t_s)] M_*/V_{cell} \quad (6d)
$$

where $e_{th}$ and $e_{nth}$ are the thermal and non-thermal energy densities, and $f_{th}$ and $f_{kn}$ are the fractions of thermal and kinetic energy deposited in the ISM. Thus, $e_{nth}$ accounts for the momentum injection by SN and $e_{th}$ for the thermal pressure part.

In the present work, we have developed a novel method to compute such quantities. The method derives $f_{th}$ and $f_{kn}$ from a detailed modelling of the subgrid blastwave evolution produced by the SN explosion. We calculate $f_{th}$ and $f_{kn}$ by evaluating the shock evolution at time $\Delta t$, the step time of the simulation$^9$.

The adopted blastwave model is based on Ostriker & McKee (1988, hereafter OM88), and it accounts for the evolution of the blast through its different evolutionary stages (energy conserving, momentum conserving, etc.). While each stage is self-similar, the passage from one stage to the next is determined by the cooling time. Thus, $f_{th}$ and $f_{kn}$ depends on the blastwave evolutionary stage. The latter, in turn depends on the gas density, cooling time, and the initial energy of the blast ($E_0 = [e_{sn}(t_s + \Delta t) - e_{sn}(t_s)] M_*$). The full model is presented in App. A.

The model details are presented in App. A. As an example, in Fig. 3, we show the energy evolution for a single SN explosion ($E_0 = 1$ foe) in a gas characterized by $n = 1 \text{ cm}^{-3}$ and $Z = 10^{-3} Z_\odot$ as a function of the time interval from the explosion $\Delta t$. We plot the total ($f = f_{th} + f_{kn}$), thermal ($f_{th}$), and kinetic ($f_{kn}$) energy fraction acquired by the gas (see eq. 6c) with solid black, dashed red and dotted blue lines, respectively. Shaded regions indicate different stages of the SN evolution, i.e. the energy conserving Sedov-Taylor (ST) stage, shell formation (SF) stage, and pressure driven snowplow (PDS). In the adopted formalism, the initial energy $E_0$ is a function of the stellar input (see Sec. 2.3 and eq. 5c), e.g.

$$
E_0 = [e_{sn}(t_s + \Delta t) - e_{sn}(t_s)] M_*. \quad \text{Full model is presented in App. A.}
$$

The underlying assumption is that the shock fronts exit the cell in $\lesssim \Delta t$. This is quite consistent because the shock is expected to be supersonic, and the sound crossing time is larger or comparable with the simulation time step $\Delta t$, dictated by the Courant-Friedrichs-Lewy conditions.

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Finally, during the PDS stage the ratio of thermal to kinetic is \( f_{\text{th}}/f_{\text{kin}} \approx 2 \) (see eqs. 6.14 in OM88).

In this particular example – a 1 foe SN exploding in a \( n = 1 \, \text{cm}^{-3} \) cell – by assuming a simulation time step of \( \Delta t \approx 10^{-2}\text{Myr} \), we find that the blastwave is in the PDS stage, and the gas receives (via eqs. 6c) a fraction of energy \( f_{\text{th}} \approx 8\% \) and \( f_{\text{kin}} \approx 16\% \) in thermal and kinetic form, respectively. During \( \Delta t \), about \( \approx 75\% \) of the initial SN energy has been either radiated away or lost to work done by the blastwave to exit the cell. The model is in broad agreement with other more specific numerical studies (e.g. Cioffi et al. 1988; Walch & Naab 2015; Martizzi et al. 2015).

### 2.4.2 Stellar winds

Stellar winds are emitted in a manner paralleling the above scheme for SNe. The energy variation can be calculated via eq. 5c, where \( \epsilon \) is directly added to some of the SPH particles, we would have

\[
\epsilon_k = 0.5 \left\langle m_b (\mathbf{v} + \Delta v \mathbf{\hat{V}}) \right\rangle V_{\text{cell}}
\]

where \( \mathbf{v} \) is the original particle velocity, the \( \langle \rangle \) operator indicates the particles sum weighted by the SPH kernel, and \( V_{\text{cell}} \) is the kernel volume. Thus, because of the kick, the increase of energy density would be

\[
\Delta \epsilon_k = \frac{0.5 \left( m_b \Delta v \mathbf{\hat{v}} + 0.5(\Delta v)^2 \right)}{V_{\text{cell}}} = 0.5 \frac{m_b (\Delta v)^2}{V_{\text{cell}}} = 0.5 \frac{m_b V_{\text{cell}}}{V_{\text{cell}}},
\]

where \( \Delta v \) can be calculated via eq. 7a, and eq. 7c can be directly cast into the AMR formalism. Additionally, because of our approximate treatment of IR-trapping (see App. B), we force energy conservation: \( f_{\text{cell}} \Delta \epsilon_k \leq \Delta t (L_{\text{ion}} + L_{\text{UV}}) \), i.e. the deposited energy must not exceed the radiative input energy. Finally, we recall here that non-thermal energy is dissipated with a time scale \( t_{\text{diss}} \), as described in the beginning of Sec. 2.

### 3 RESULTS

At \( z = 6 \) (\( t \approx 920 \text{Myr} \)), the simulated zoom-in region contains a group of 15 DM haloes that host galaxies. We target the most massive halo (\( M_h = 1.8 \times 10^{13} M_\odot \)) that hosts “Dahlia”, which is a galaxy characterized by a stellar mass of \( M_* = 1.6 \times 10^{10} M_\odot \), therefore representative of a typical LBG galaxy at that epoch. Dahlia has 14 satellites located within \( \approx 100 \text{ kpc} \) from its centre. The six largest ones have a DM mass in the range \( M_h = 2.5 \times 10^{10} M_\odot - 1.2 \times 10^{10} M_\odot \), and they host stars with total mass \( M_\star \lesssim 10^9 M_\odot \). Additionally, there are eight smaller satellites (\( M_h \approx 10^9 M_\odot \)), with \( M_\star \approx 10^8 M_\odot \).

#### 3.1 Overview

We start by looking at the overall properties of Dahlia on decreasing scales. In the following we refer to Fig. 4, which shows the simulated density \( n \), temperature \( T \), total (thermal+kinetic) pressure \( P \), and metallicity \( Z \) maps at \( z = 6 \).

#### 3.1.1 Environment (scale \( \approx 160 \text{ kpc} \))

Dahlia sits at the centre of a cosmic web knot and accretes mass from the intergalactic medium (IGM) mainly via 3 filaments of length \( \approx 100 \text{ kpc} \), confirming previous findings (Dekel et al. 2009). These overdense filaments (\( n \approx 10^{-2} \text{cm}^{-3} \)) are slightly colder (\( T \approx 10^{3.5} \text{K} \)) than the IGM.

---

\(^{10}\) In eq. 7c, when going from the first to the second line, note that first terms gives a null contribution, as \( \mathbf{v} \) is ordered motion, while the kicks are randomly oriented via \( \mathbf{v} \) and that, by definition, \( \langle \mathbf{P}_{\text{kick}} \rangle = 1 \). 

\(^{11}\) Most of the maps of this paper are obtained with a customized version of PYNMES (Labadens et al. 2012), a visualization software that implements optimized techniques for the AMR grid of RAMSES.
Figure 4. (Caption next page.)
\[ (T) \approx 10^{6.5} \text{K} \] as a consequence of their shorter radiative cooling time \( (\tau_{\text{cool}} \propto n^{-1}) \). Along these cold streams, pockets of shock-heated \( (T \gtrsim 10^4 \text{K}) \) gas produced by both structure formation and feedback (SN and winds) are visible.

The galaxy locations can be pinpointed from the metallicity map, showing a dozen of metal-polluted regions. The size of the metal bubbles ranges from \( \approx 20 \text{kpc} \) in the case of Dahlia to a few kpc for the satellites. Bubble sizes increase with the total stellar mass (see P14, in particular Fig. 13), and age of the galaxy stellar population.

On these scales, the pressure is dominated by the thermal component \( (P \approx P_{\text{th}} \approx 10^4 \text{K} \text{cm}^{-3}) \); higher values of pressure, associated to non-thermal feedback effects (e.g. gas bulk motion), are confined around star forming regions, again traced by the metallicity distribution.

### 3.1.2 Circumgalactic medium (scale \( \approx 50 \text{kpc} \))

To investigate the circumgalactic medium (CGM), we zoom in a region within \( \sim 3r_{\text{vir}} = 47.5 \text{kpc} \) from Dahlia’s centre. On these scales, we can appreciate the presence of several Dahlia’s satellites, i.e. extended (few kpc) structures that are \( \sim 100 \) times denser than the filament in which they reside. Two of these density structures are particularly noticeable. These are located at a distance of \( \sim 10 \text{kpc} \) from the centre in the upper left and lower left part of the map, respectively. By looking at the metallicity distribution, we find that both satellites reside within their own metal bubble, which is separated from Dahlia’s one. This clearly indicates an in-situ star formation activity.

Additionally, the density map shows about 20 smaller \( (\sim 10 - 100 \text{pc}) \) overdense clumps \( (n \gtrsim 10^2 \text{cm}^{-3}) \). The ones within Dahlia’s metal bubble are enriched to \( Z \gtrsim Z_\odot \). This high Z value is indicative of in-situ self-pollution, which possibly follows an initial pre-enrichment phase from Dahlia. Clumps outside Dahlia metal bubble have on average a higher density \( (n \sim 10^3 \text{cm}^{-3}) \). Since these clumps are unpolluted, they have not yet formed stars, as the effective density threshold for star formation is \( \sim 25/(Z/Z_\odot) \text{cm}^{-3} \) (see eq. 3b and Sec. 2.2). Such clumps represent molecular cloud complexes caught in the act of condensing as the gas streams through the CGM (Ceverino et al. 2016). Such clumps have gas mass in the range \( 10^5 - 10^6 \text{M}_\odot \), and are not DM-confined, as the DM density field is flat on their location.

Star forming regions are surrounded by an envelope of hot \( (T \sim 10^{5.5} \text{K}) \), diffuse \( (n \gtrsim 10^{-2} \text{cm}^{-3}) \) and mildly enriched \( (Z \sim 10^{-2} Z_\odot) \) gas produced by SN explosions and winds. In the centre of star forming regions, instead, the gas can cool very rapidly due to the high densities/metallicities. Nevertheless, these regions are highly pressurized due to bulk motions mostly driven by radiation pressure (see Fig. 8).

### 3.1.3 ISM (scale \( \approx 10 \text{kpc} \))

The structure of Dahlia’s ISM emerges once we zoom in a region \( \sim 0.5r_{\text{vir}} \) from its centre. In the inner region \( (\sim 2 \text{kpc}) \), a counterclockwise disk spiral pattern is visible, since the field of view is perpendicular to the rotation plane of the galaxy (see Gallerani et al. 2016 for the analysis of the velocity field of Dahlia). The presence of disks in these early systems has already been suggested by other studies. For example, Feng et al. (2015) show that already at \( z \sim 8 \) nearly 70% of galaxies with \( M_* \sim 10^{10} \text{M}_\odot \) have disks (see also Sec. 3.3).

The spiral central region and the spiral arms are dense \( (n \sim 10^3 \text{cm}^{-3}) \) and cold \( (T \sim 10^4 \text{K}) \), and the active SF produces a large in-situ enrichment \( (Z \gtrsim Z_\odot) \). Winds and shocks from SN have no effect in the inner part of the galaxy, because of the high density and short cooling time of the gas; this implies that metals remain confined within \( \sim 2 \text{kpc} \). Within spiral arms radiation pressure induced bulk motions largely dominate the total pressure, which reaches values as high as \( P \gtrsim 10^6 \text{K} \text{cm}^{-3} \). The imprint of SN shocks is evident in the temperature map in regions with \( T \gtrsim 10^7 \text{K} \). Shock driven outflows originated in spiral arms travel outward in the CGM, eventually reaching the IGM if outflow velocities exceed the escape velocity \( (\sim 100 \text{km s}^{-1}) \), see Fig. 4 in Gallerani et al. 2016.

Outflows are either preferentially aligned with the galaxy rotation axis, or they start at the edge of the disk. However, when spherically averaged, infall and outflow rates are nearly equal \((\sim 30 \text{M}_\odot/\text{yr} \text{ at } z \sim 6, \text{ Gallerani et al. 2016}), \text{ and the system seems to self-regulate (see also Dekel & Mandelker 2014).}

Outside the disk, clumps with density \( n \sim 10^3 \text{cm}^{-3} \) are also present and are actively producing stars. These isolated star forming MCs are located at a distance \( \gtrsim 2 \text{kpc} \) from the centre, and show up as spots of high pressure \( (P \gtrsim 10^5 \text{K} \text{cm}^{-3}) \); some of these MCs are completely disrupted by internal feedback and they can be recognized by the low metallicity \( (Z \sim 10^{-2} Z_\odot) \); this is consistent with the outcome of numerical simulations of multiple SN explosions in single MC (e.g. Körtgen et al. 2016).

### 3.1.4 Radial profiles

Fig. 5 shows spherically averaged density, metallicity, and \( \text{H}_2 \) density profiles for the gas. The density profile rapidly decreases from \( n \sim 30 \text{cm}^{-3} \) at \( r \sim 0 \) to \( n \sim 0.1 \text{cm}^{-3} \) at \( r \sim 6 \text{kpc}(\sim 0.5r_{\text{vir}}) \), and then flattens at larger distances. Such profile is consistent with the average profile of \( z = 4 \) galaxies presented in P14. There we claimed that the density profile is universal once rescaled to the halo virial radius (see also Liang et al. 2016). Superposed to the mean density profile, local peaks are clearly visible: they result from individual clumps/satellites, as discussed above.

The central metallicity is close to the solar value, but by \( r \sim 12 \text{kpc} \sim r_{\text{vir}} \) it has already dropped to...
$Z = Z_{\text{floor}}$. Within $0 \lesssim r/\text{kpc} \lesssim 6$, the metallicity gradient closely tracks the density profile, while for $6 \lesssim r/\text{kpc} \lesssim 15$ the decrease is steeper. Pallottini et al. (2014b) find that the metallicity profile is not universal, however it usually extend up to few virial radii, as for Dahlia; further insights can be obtained by analyzing the $n-Z$ relation (Sec. 3.2.3).

In Fig. 5 we note that the $Z$ gradient found in Dahlia at $z = 6$ is slightly steeper than the one inferred from observations of $z \sim 3$ galaxies: i.e. we find $\Delta Z/\Delta r \sim -0.1 \text{ dex}/\text{kpc}$ while the observed ones are $\sim 0 \text{ dex}/\text{kpc}$ (Wuyts et al. 2016) and $\sim +0.1 \text{ dex}/\text{kpc}$ (Troncoso et al. 2014). This suggests that the metallicity profile evolve with cosmic time and that the flattening is likely caused by stellar feedback, which in our Dahlia may occur in the following Gyr of the evolution. However, to prove such claim we should evolve the simulations to $z \sim 3$.

The $H_2$ profile is spiky, and each peak marks the presence of a distinct SF region\(^\text{12}\). In Dahlia $H_2$ is mainly concentrated within $r \lesssim 0.5 \text{kpc}$ and it is distributed in the disk-like structure seen in Fig. 4 (see Sec. 3.3). The location of the other peaks correspond to the satellites, which are mostly co-located with metallicity peaks. With increasing metallicity, in fact, lower densities are needed to form $H_2$ (eq. 3b).

### 3.2 Star formation and feedback history

We analyze the SF history of Dahlia and its major satellites by plotting in Fig. 6 the cumulative stellar mass ($M_*$) and star formation rate ($SFR$) vs. time\(^\text{13}\).

For the whole galaxy sample, the time averaged ($\pm \text{r.m.s.}$) specific star formation is $\langle SFR \rangle = 16.6 \pm 32.8 \text{ Gyr}^{-1}$. This mean value is comparable to that obtained by previous simulations of high-$z$ galaxies (Wise et al. 2012) and broadly in agreement with $z \sim 7$ observations (Stark et al. 2013). At early times the $sSFR$ reaches a maximum of $\sim 100 \text{ Gyr}^{-1}$, while a minimum of $3.0 \text{ Gyr}^{-1}$ is found during the late time evolution. Both the large $sSFR$ range and maximum at early times are consistent with simulations by Shen et al. (2014). At late times, the $SFR$ is in agreement with analytical calculation (Behroozi et al. 2013), and with $z = 7$ observations (Gonzalez et al. 2010), although we note that Dahlia has a larger stellar mass with respect to the galaxies in the sample ($M_* \approx 5 \times 10^{10} M_\odot$).

At all times, Dahlia dominates both the stellar mass and star formation rate, whose mean value is $\langle SFR \rangle \approx (35.3 \pm 32.7) M_\odot/\text{yr}$. Its stellar mass grows rapidly, and it reaches $M_* \approx 10^8 M_\odot$ by $t \approx 400 \text{ Myr}$, i.e. after $\approx 120 \text{ Myr}$ from the first star formation event. Such rapid mass build-up is due to merger-induced SF, that plays a major role at high-$z$ (Poole et al. 2016; Behroozi et al. 2013; Salvadori et al. 2010). The $SFR$ is roughly constant from $z \sim 11$ to $z \sim 8.5$ and reaches a maximum of $\approx 130 M_\odot/\text{yr}$ at $z \sim 6.7$. With respect to observations of $z \lesssim 6$ LBG galaxies (e.g. Stanway et al. 2003; Stark et al. 2009) the $SFR$ and $M_*$ of Dahlia are above the mean values, but still consistent within one sigma. Additionally, the combination of $SFR$, $M_*$, and $Z$, for Dahlia are compatible with the fundamental mass metallicity relation observed in local galaxies (Mannucci et al. 2010).

The total stellar mass in satellites is $M_* \approx 10^8 M_\odot$. Typically, SF starts with a burst, generating $\sim 10^{-8} M_\odot$ of stars during the first $\approx 20 \text{ Myr}$. Then the $SFR$ exponentially declines and becomes intermittent with a bursty duty cycle of $\sim 100 \text{ Myr}$. This process can be explained as follows. As an halo forms, at its centre the density of the gas slowly rises. When the density is higher than the critical density of $H_2$ formation (eq. 3b), the gas in the inner region is converted into stars in few free-fall times. Then feedback, and in particularly coherent SN explosions ($t \gtrsim 10 \text{ Myr}$, see Fig. 2), quenches $SFR$, and the star formation activity becomes self-regulated. As mergers supply fresh gas, the $SFR$ suddenly goes out of equilibrium and becomes bursty again. Note that self-regulation is possible only for major satellites, since smaller ones ($M_* \lesssim 10^8 M_\odot$) cannot retain a large fraction of their gas following feedback events due to their shallow potential wells (see P14).

Note that the duty cycle and the amplitude of the burst are fairly in agreement with observations of $M_* \approx 10^8 − 10^{10} M_\odot$ galaxies at $z \lesssim 0.3$ (Kaufmann 2014). Furthermore, in our satellites we find that the typical behavior of the burst phases – starburst - quiescent - post-starburst – is qualitatively similar to what found by Read et al. (2016b), that simulate the evolution of a $M_* \approx 10^8 M_\odot$ galaxy for $\approx 1 \text{ Gyr}$ (see also Teyssier et al. 2013; Read et al. 2016a for further specific studies on the bursty nature of this kind of galaxies).

Since individual galaxies are defined as group of star particles in the same DM halo at $z = 6$, the SF history accounts for the sum of all the stars that formed in different progenitors of the considered halo. For comparison, in the right panel of Fig. 6 we plot the $SFR$ of individual halos defined by their merger history at $z = 8.7$. Galaxies with active SF at $300 − 550 \text{ Myr}$ merge into Dahlia at a later...
The internal structure of $z \simeq 6$ galaxies

Figure 6. **Left**: Dahlia and satellites cumulative stellar masses ($M_\star$, upper left panel) and star formation rates (SFR, lower left panel) as a function of cosmic time ($t$). For each galaxy, individual $M_\star$ and SFR are plotted with a solid line, coloured accordingly to the total dark matter mass ($M_{\text{h}}$) of the host halo at $z = 6$. For both $M_\star$ and SFR, Dahlia’s tracks are plotted with a blue line, and the totals (Dahlia+satellites) are in black. **Right**: SFR as a function of cosmic time, with individual galaxies defined by the merger history up to $z \simeq 8.5$. Note the different $M_{\text{h}}$ colourbar scale with respect to the left panel.

Figure 7. Effective star formation efficiency ($\zeta_{\text{sf}} f_{\text{H}_2}$) vs density ($n$) at $z = 6$. The distribution is H$_2$ mass weighted; we consider gas within $3 r_{\text{vir}} = 47.5$ kpc from Dahlia centre.

3.2.1 Star formation efficiency

$\zeta_{\text{sf}} f_{\text{H}_2}$ represents the quantity of gas converted in stars within a free-fall time (see eq. 1). In Fig. 7 we plot the effective star formation efficiency ($\zeta_{\text{sf}} f_{\text{H}_2}$) as a function of gas density, weighted by the H$_2$ mass fraction at $z = 6$. Most of the H$_2$ is contained in the range $n = 10 - 100$ cm$^{-3}$, and the effective efficiency $\zeta_{\text{sf}} f_{\text{H}_2}$ varies from $10^{-3}$ to $10^{-1}$. Since $\zeta_{\text{sf}} = \text{const.} = 0.1$, the spread is purely due to the dependence of $f_{\text{H}_2}$ on density and metallicity (see Fig. 1). Note that by construction $\zeta_{\text{sf}} f_{\text{H}_2} \leq 0.1$, and the plot does not show values very close to such limit, since gas with higher effective efficiency is converted into stars within a few free-fall times (eq. 1b).

Interestingly, our H$_2$-based star formation criterion is reminiscent of a density threshold one, as below $n \simeq 3$ cm$^{-3}$ the efficiency drops abruptly (eqs. 3). However, an important difference remains, i.e. in the present model at any given density the efficiency varies considerably as a result of the metallicity dependence. The relation between efficiency and density is also similar to that found by Semenov et al. (2016) (S16). This is striking because these authors use a star formation efficiency that depends on the turbulent velocity dispersion of the gas, with no notion of the local metallicity. This comparison is discussed further in Sec. 4.

3.2.2 Feedback energy deposition

As discussed in Sec. 2.4, only a small fraction of the available energy produced by stars can couple to the gas. During the simulation, we find that the time average efficiency of the conversion is $f \sim 0.1\%$, regardless of the feedback type. These low efficiencies imply that energy is mostly dissipated within MCs where the stars reside and produce it. For SN and winds, such small efficiency is a consequence of the short cooling times in MCs (see also App. A). For radiation pressure the efficiency is limited by the relatively small dust optical depths (see also App. B).

Note that, typically in simulations (e.g. Wise et al. 2012; Agertz et al. 2013), energy from stars is directly deposited in the gas, and then dissipation (mostly by radiative losses) occurs during the hydrodynamical time step. Within our scheme, instead, the deposited energy is already dissipated within high density cells, where cooling is important. Nevertheless, this does not appear to determine major differences.
in, e.g., SFR history and ISM thermodynamics, as discussed in Sec. 3.2.3.

In Fig. 8 we plot the energy deposition rate in the gas by various feedback processes as a function of time. Most evidently, radiation dominates the energy budget at all times: $E_{\text{rad}} \simeq 10^7 E_{SN} \simeq 10^7 E_w$. The ratios of these energy rates somewhat reflect the stellar inputs shown in Fig. 5, although this is not a trivial finding, given that the interplay among different feedback types is a highly non-linear process.

As expected, the energy deposition rate behaves as $\dot{E} \propto SFR^q$, with $q \gtrsim 1$, apart from fluctuations and jumps as the one at $t \simeq 500\,\text{Myr}$. The scaling can be understood by simple dimensional arguments. Assume that most of the energy is deposited by radiation pressure. In the optically thick limit, we can combine eqs. 7c and 7a to write $\dot{E}_{\text{rad}} \Delta t \simeq (L_{\text{rad}} \Delta t) / (M_g c^2)$, where $M_g$ is the gas mass accelerated by radiation, and we neglect ionizing radiation. Then, using 5e, we can write $\dot{E}_{\text{rad}} \propto SFR (M_g / M_*)$. Initially, $M_g \simeq M_*$, thus $\dot{E}_{\text{rad}} \propto SFR$. Once the gas mass is expelled from the star forming region or converted into stars, $M_g \ll M_*$ Thus the deposition rate increases faster than the SFR and it is very sensitive to the amount of gas mass around the sources.

The previous argument holds until the gas remains optically thick. This is warranted by AGB metal/dust production which becomes important after for stellar ages $t_* \sim 100\,\text{Myr}$ (see Fig. 2). When combined with the parallel increase of UV photons by the same sources, it is easy to interpret the rapid increase of the radiative feedback efficiency at $t \simeq 500\,\text{Myr}$, i.e. after $\sim 200\,\text{Myr}$ from the first star formation events in Dahlia. We checked this interpretation by looking at the IR-trapping recorded on the fly during the simulation. We find that on average $f_{ir} \simeq 10^{-2}$ for $t \lesssim 500\,\text{Myr}$, and $f_{ir} \simeq 0.1$ at later times, thus confirming our hypothesis.

The energy deposition rates for different feedback types are highly correlated in time (Pearson coefficients $\gtrsim 0.7$). This is partially due to the fact that the same stellar population inputs wind, radiation and supernova energy in the gas. Additionally, as we have just seen for the case of AGB star, different types of feedback are mutually dependent. For example, radiation pressure is more effective when the gas is metal and dust enriched by SN and AGB stars; winds and SN can more efficiently couple with low density gas (longer cooling time).

Note that short and intense peaks in energy deposition rate correspond to the complete disruption of multiple MCs. This occurs following strong SF events in small satellites ($M_\bullet \sim 10^7 M_\odot$) that cannot retain the gas and sustain a continuous star formation activity.

Finally, we remind that, when compared with observational/analytical constraints, the SFR and $M_*$ of Dahlia are higher then the mean, but still consistent within one sigma. We caution that this might imply a somewhat weak feedback prescription.

### 3.2.3 Feedback effects on ISM thermodynamics

Feedback leaves clear imprints in the ISM thermodynamics. For convenience, we classify ISM phases according to their density: we define the gas to be in the rarefied, diffuse, and dense phase if $n < 0.1\,\text{cm}^{-3}$, $0.1 \leq n < 0.1 \times 10^2\,\text{cm}^{-3}$, and $n > 10^2\,\text{cm}^{-3}$, respectively.

We focus at $z = 6$ and consider the gas in a region within $\sim 47.5\,\text{kpc}$ (3$r_{vir}$) from Dahlia’s centre, essentially the scale of the CGM described in Sec. 3.1.2. This region contains a total gas mass of $1.3 \times 10^{10} M_\odot$, and metal mass of $4.2 \times 10^8 M_\odot$ (additional data in Tab. 2).

Fig. 9 shows the Equation of State (EOS, or phase diagram) of the gas. The fraction of gas in the rarefied, diffuse and dense phases is 44%, 34% and 22%; these phases contain 5%, 25% and 70% of the metals, respectively. Thus, while the gas mass is preferentially located in the lower density phases, metals are mostly found in dense gas, i.e. star forming regions/MC. Additionally only $\sim 30\%$ of the considered

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14 Compared to the definitions used in Klessen & Glover (2014), the rarefied corresponds to the warm and hot ionized medium, the diffuse phase to the cold and warm neutral medium and the dense phase to the molecular gas.

| mass   | rarefied | diffuse | dense |
|--------|----------|---------|-------|
| Gas    | $1.3 \times 10^{10} M_\odot$ | 44%    | 34%   | 22%   |
| Metals | $4.2 \times 10^{9} M_\odot$  | 5%     | 25%   | 70%   |
| H$_2$  | $3.6 \times 10^{8} M_\odot$  | 0%     | 1%    | 99%   |
| C II   | $2.2 \times 10^{7} M_\odot$  | 4%     | 22%   | 74%   |

Table 2. Summary of the gas masses for total, metal, C II, and H$_2$ within $\sim 47.5\,\text{kpc}$ (3$r_{vir}$) from Dahlia center. In the table, we report also the fraction that is contained in different gas phases: rarefied ($\log(n/\text{cm}^{-3}) \leq -1$), diffuse ($-1 < \log(n/\text{cm}^{-3}) \leq 1$) and dense ($\log(n/\text{cm}^{-3}) > 1$). Discussion about gas and metal mass is found in Sec. 3.2.3; analysis of H$_2$ and C II is in Sec. 3.3 (see also App. C for C II calculation).

Figure 8. Rate of energy deposition in the gas, $dE/dt$, by feedback processes as a function of cosmic time. Different contributions (SN, wind and radiation) are plotted with a different colour, and we additionally distinguish between the kinetic (thick lines) and thermal (thin lines) energy variation. By definition, radiation pressure has no thermal contribution. Note the jump at $t \sim 500\,\text{Myr}$ due to the onset of radiation pressure by AGB stars. The upper axis indicates the corresponding redshift.
The internal structure of \( z \simeq 6 \) galaxies

Figure 9. Equation of State of the gas within \( \simeq 47.5 \) kpc (3 \( r_{\text{vir}} \)) from Dahlia centre at \( z = 6 \). Each EOS consists in a mass- or metal-weighted probability distribution function (PDF) as specified by the colourbar. We plot the PDF in the \( n-T \) plane (upper left panel), the metal-mass weighted PDF in the \( n-T \) plane (lower left panel), and mass-weighted relation between gas \( n \) and \( Z \) (lower left right). Mean relations and r.m.s. dispersions are overplotted with solid black and dashed lines, respectively. In the upper horizontal axis of each panel, we indicate the overdensity \( \Delta \) corresponding to \( n \). The density range of rarefied, diffuse and dense phases used in the text are indicated. For the panels on the left, the rarefied gas is additionally divided in photo-ionized (\( T < 10^4.5 \) K) and shock-heated (\( T \geq 10^4.5 \) K). See Tab. 2 for a summary of the total values.

volume shows \( Z > 10^{-3}Z_{\odot} = Z_{\text{floor}} \), i.e. it has been polluted by stars in the simulation. We note that the EOS in the \( n-T \) plane is fairly consistent with the one found in other high-\( z \) galaxy simulations (e.g. see Fig. 5 in Wise et al. 2012). Comparison between the EOS in the \( n-T \) and \( n-P \) plane highlights the relative importance of different feedback types.

The rarefied gas is characterized by long cooling times. Thus, once engulfed by shocks, such phase becomes mildly enriched \( \langle Z \rangle \sim 10^{-2}Z_{\odot} \) and remains hot (\( T \sim 10^6 \) K). The enriched rarefied gas preferentially populates the \( n \sim 10^{-3} \) cm\(^{-3} \) and \( T \sim 10^6 \) K region of the phase diagram. However, part of the rarefied gas has \( T \sim 10^4 \) K. This gas component has a temperature set by the equilibrium between adiabatic cooling and the photo-heating by the UV background; it feeds the accretion onto Dahlia, but it is not affected by stellar feedback. As such it is not central in the present analysis.

The dense gas is mostly unaffected by shocks and it is concentrated in the disk. Typically, such gas has \( n \sim 10^2 \) cm\(^{-3} \) and \( T \sim 10^2 \) K, thus a thermal pressure \( P_{\text{th}}/k \sim 10^4 \) cm\(^{-3} \) K is expected. However, the total gas pressure is \( P/k \sim 10^7 \) cm\(^{-3} \) K (see the \( P-n \) EOS). The extra contribution is provided in kinetic form by radiation pressure, thanks to the strong coupling with the gas allowed by the high optical depth of this phase. This leads to the important implication that the central structure of Dahlia is radiation-supported (see also Sec. 3.3).

The diffuse gas acts as an interface between the dense disk gas and the rarefied gas envelope. Diffuse gas is found both in hot (\( T \sim 10^6 \) K) and cold (\( T \sim 10^4 \) K) states. The cold part has a sufficiently high mean metallicity, \( Z \sim 0.1Z_{\odot} \), to allow an efficient cooling of the gas. This is highlighted by the metal-weighted EOS, where we can see that most of the metals present in the diffuse phase are cold.

Note that the phase diagram also shows evidence for the classical 2-phase medium shape for pressures around \( P/k \sim 10^5 \) cm\(^{-3} \) K, while at higher (and lower) pressures only one stable phase is allowed; nevertheless, at any given pressure a
of \( \sim 15 \, M_\odot \, \text{yr}^{-1} \, \text{kpc}^{-2} \), i.e. more than 1000 times the Milky Way value. Fragmentation of the disk is relatively weak (cfr. Mayer et al. 2016), as indicated by smooth surface density map, and also paralleling the flat metallicity profile in the inner \( z \approx \text{kpc} \). For the fragmentation of the \( \text{H}_2 \) component, we caution that this result has been obtained assuming a uniform UV interstellar field; stronger fragmentation in the \( \text{H}_2 \) distribution may occur when accounting for local radiation sources: Lyman-Werner photons from these sources might in fact locally dissociate the \( \text{H}_2 \) by generating pockets of \( \text{H}_1 \) in the distribution.

While most of the \( \text{H}_2 \) gas resides in the disk, we can clearly distinguish 3 clumps of molecular gas both in the face-on and edge-on maps. These clumps are located few kpc away from the centre, and are characterized by sizes of \( \sim 150 \, \text{pc} \) and \( M_{\text{HI}} \sim 5 \times 10^6 M_\odot \). Such clumps are Jeans-unstable and form stars as they infall and stream through the CGM, as it can be appreciated by comparing \( \text{H}_2 \) and stellar mass profiles (Fig. 11). The stellar mass profiles also highlight the presence of 3 stellar clumps at \( r \sim 1 \, \text{kpc} \) with no associated \( \text{H}_2 \). These “older” clumps share the same nature of the previous ones, but the H2 has been already consumed and/or dispersed by the star formation activity that produced the stars present at \( z = 6 \).

### 3.3 Additional ISM properties

We conclude our analysis by inspecting the distribution of two key ISM species, molecular hydrogen and \( \text{C} \, \text{II} \), along with the expected surface brightness of the corresponding 158\,\mu m \( \text{C} \, \text{II} \) line. The surface maps of these quantities in Dahlia (\( z = 6 \)) are shown in Fig. 10 for the face-on and edge-on view cases. For reference, in Fig. 11 we additionally plot the radially-averaged profiles of the same quantities, and in Tab. 3 we give their typical radial scales.

#### 3.3.1 Molecular Hydrogen

Dahlia has a total \( \text{H}_2 \) mass \( M_{\text{H}_2} \sim 3.6 \times 10^8 M_\odot \), that is mainly concentrated in a disk-like structure of radius \( \sim 0.6 \, \text{kpc} \) and scale height \( \sim 200 \, \text{pc} \), with a sharp cut off beyond these scales\(^{15}\). The disk has mean surface density \( \langle \Sigma_{\text{H}_2} \rangle \sim 10^{2.5} M_\odot \, \text{kpc}^{-2} \), that is approximately constant with radius and presents perturbed spiral arms along which the density is enhanced by a factor \( \sim 3 \). The spiral arms are less pronounced than in a more massive, MW-like galaxy (see S16 and Ceverino et al. 2015). This trend with mass has already been pointed out by Ceverino et al. (2010).

The disk is composed by dense \(( n \sim 25 \, \text{cm}^{-3}) \), enriched \( Z \sim 0.5 Z_\odot \), radiation-pressure supported gas, as already discussed. It is fed by frequent mergers driving fresh gas to the centre, and supports a star formation rate per unit area

| \[ r_{1/2}/\text{kpc} \] | \[ \text{face-on} \] | \[ \text{edge-on} \] | \[ \text{value at } r_{1/2} \] |
|-----------------|-----------------|-----------------|------------------|
| \( \text{H}_2 \) | 0.59 | 0.36 | \( 10^2.23 M_\odot \) |
| \( \text{C} \, \text{II} \) | 0.64 | 0.38 | \( 10^5.14 M_\odot \) |
| stars | 0.37 | 0.23 | \( 10^7.89 M_\odot \) |
| \( \text{[C} \, \text{II}] \) | 0.60 | 0.36 | \( 10^7.25 L_\odot \) |

#### 3.3.2 Singly ionized Carbon

The \( \text{C} \, \text{II} \) abundance is calculated by post-processing the simulation outputs with the photoionization code CLOUDY (Ferland et al. 2013, and see App. C). The result is shown in Fig. 10. Dahlia contains a \( \text{C} \, \text{II} \) mass of \( M_{\text{CII}} = 2.2 \times 10^8 M_\odot \), accounting for \( \sim 50\% \) of the total metals produced. About 74\% of the \( \text{C} \, \text{II} \) mass is located in the dense phase, 22\% in the diffuse phase, 4\% in the rarefied phase. Note that the \( \text{C} \, \text{II} \) mass phase distribution differs only for \( \lesssim 10\% \) from the \( Z \) distribution (see Tab. 2). The difference arises because shock-heated gas can be collisionally excited to higher ionization states. Thus, to first order, we expect the \( \text{C} \, \text{II} \) spatial distribution to follow the metallicity one.

The face-on \( \text{C} \, \text{II} \) surface density has a central maximum \( (\Sigma_{\text{CII}} \sim 10^9 M_\odot \, \text{kpc}^{-2}) \), it gradually decreases to up to \( \sim 1.2 \, \text{kpc} \), and drastically drops to \( \Sigma_{\text{CII}} \lesssim 10^7 M_\odot \, \text{kpc}^{-2} \) beyond that radius (see also Fig. 11). Thus, most of the \( \text{C} \, \text{II} \) is located into the disk, but a more extended envelope containing a sizable fraction of mass exists. On top of this smooth distribution, there are \( \text{C} \, \text{II} \) enhancements corresponding to the \( \text{H}_2 \) clumps described above.

The \( \text{C} \, \text{II} \) profile is similar for edge-on and face-on case. However, the edge-on has a higher \( \text{C} \, \text{II} \) central density and a steeper slope. While the higher central value is obviously due to the larger column density encountered along the disk, the sharp drop is related to metal transport. As most of the star formation activity is located in the disk, metals above it can be only brought by outflows which become progressively weaker with distance. Metal outflows originating from the centre are preferentially aligned with the rotation axis, and the pollution region starting from the edge is stretched by the disk rotation and by tidal interaction with satellites.
3.3.3 Emission from singly ionized carbon

We finally compute the expected [C\textsc{ii}] line emission using the same prescriptions of Vallini et al. (2013, 2015), as detailed in App. C. Note that for the present work we assume uniform UV interstellar radiation. This approximation is valid in the MW, where variations around the mean field value are limited to a factor of 3. The results are plotted in Fig. 10.

Within 1 kpc from the centre the [C\textsc{ii}] emission structure closely follows the C\textsc{ii} distribution, and we find $S_{\text{CII}}/L_{\odot} \approx 200 \Sigma_{\text{CII}}/M_{\odot}$. At larger radii the [C\textsc{ii}] surface brightness suddenly drops, although the peaks associated with H$_2$ clumps are preserved. This result holds both for the face-on and edge-on cases.

Such behavior can be understood as follows. Take a typical MC with $n = 10^2$ cm$^{-3}$, $Z = Z_{\odot}$, and total mass $M$. Its [C\textsc{ii}] luminosity is $L_{\text{CII}}/L_{\odot} \approx 0.1 (M/M_{\odot})$ (Vallini et al. 2016; Goicoechea et al. 2015). Also, the [C\textsc{ii}] emission is $\propto Z n$ for $n \lesssim 10^3$, i.e. the critical density for C\textsc{ii} collisional excitation by H atoms (Vallini et al. 2013). Then,

$$L_{\text{CII}} \approx 0.1 \left( \frac{n}{100 \text{ cm}^{-3}} \right) \left( \frac{Z}{Z_{\odot}} \right) \left( \frac{M}{M_{\odot}} \right) L_{\odot}. \quad (8)$$

In the central kpc, where $n \approx 10^2$ cm$^{-3}$ and $Z \approx Z_{\odot}$, the luminosity depends only on the molecular mass contained in the disk, and the same holds even for H$_2$ clumps outside the disk. The envelope is instead more diffuse ($n \approx 100$ cm$^{-3}$) and only mildly enriched ($Z \lesssim 10^{-3} Z_{\odot}$). As a result, its [C\textsc{ii}] luminosity per unit mass is lower.

The emission from this diffuse component is further suppressed by the CMB (da Cunha et al. 2013; Pallottini et al. 2015; Vallini et al. 2015). Namely, for gas with $n \lesssim 0.1$ cm$^{-3}$, the upper levels of the [C\textsc{ii}] transition cannot be efficiently populated through collisions, thus the spin temperature of the transition approaches the CMB one, and to a first order the gas cannot be observed in emission.

In summary, $\approx 95\%$ of Dahlia [C\textsc{ii}] emission comes from dense gas located in the H$_2$ disk. Indeed, the [C\textsc{ii}] half light radius coincides with the H$_2$ half mass radius, i.e. 0.59 kpc (0.36 kpc) in the face-on (edge-on) case (see also Tab. 3). Within such radius, the molecular gas has a mass $M_{\text{H}_2} \approx 1.69 \times 10^8 M_{\odot}$ and the luminosity is $L_{\text{CII}} \approx 1.78 \times 10^5 L_{\odot}$, i.e. with a [C\textsc{ii}]-H$_2$ scaling ratio consistent within 15% from the simple estimate in eq. 8.

Dahlia has a total [C\textsc{ii}] luminosity $L_{\text{CII}} \approx 3.5 \times 10^7 L_{\odot}$; this is fainter than expected on the basis of the local [C\textsc{ii}]-SFR relation ($L_{\text{CII}} \approx 10^8 - 10^9 L_{\odot}$; i.e. De Looze et al. 2014). However, at high-$z$, such relation seems to hold only for a small subset of the observed galaxies (Capak et al. i.e. 2015; Willott et al. i.e. 2015). The majority of the observed galax-
ies show a strong [C II]-SFR deficit, when considering both
detections (e.g. BDF3299, A383-5.1 Maiolino et al. 2015;
Knudsen et al. 2016) and upper limits (e.g. Himiko, IOK1,
MS0451-H Ouchi et al. 2013; Ota et al. 2014; Knudsen et al.
2016).

For Dahlia, the [C II]-SFR deficit depends on multiple
factors. The main contribution from [C II] emission is in the
H2 disk, that on average has \( Z \approx 0.5 Z_\odot \), i.e. slightly lower
then solar. Additionally, the gas in the disk is efficiently con-
verted in stars \( (\Sigma_{\text{em}} \approx 100 M_\odot/yr) \) and has \( n \approx 25 \times 10^{-3} \),
thus the [C II] emission is hindered (eq. 8). Finally, there is
a marginal contribution to [C II] from the diffuse and rared-
ified phase; \( \approx 30\% \) of C II is locked in a the low density
and metallicity gas that gives a negligible contribution to
[C II] emission, particularly because of CMB suppression.

4 SUMMARY AND DISCUSSION

With the aim of characterizing the internal properties of
high-z galaxies, we have performed an AMR zoom-in sim-
ulation of “Dahlia”, a z \( \approx 6 \) galaxy with a stellar mass of
\( M_* = 1.6 \times 10^{10} M_\odot \), therefore representative of LBGs at
that epoch. We follow the zoom-in region with a gas mass
resolution of \( 10^9 M_\odot \) and a spatial resolution of 30 pc.

The simulation contains a rich set of physical processes.
We use a star formation prescription based on a H2 depen-
dent Schmidt-Kennicutt relation. The H2 abundance is com-
puted from the KTM09 model (Fig. 1). Using stellar evolu-
tionary models (Bertelli et al. 1994; Leitherer et al. 1999),
we include chemical, radiative and mechanical energy in-
puts, accounting for their time evolution and metallicity de-
pendence on the stellar population properties (Fig. 2). We
include feedback from SN, winds and radiation pressure with
a novel, physically motivated coupling scheme between gas
and stars. We also compute [C II] abundance and the 158\( \mu \)m
[C II] emission, by post-processing the outputs with CLOUDY
(Ferland et al. 2013), and a FIR emission model drawn from
radiative transfer numerical simulations (Vallini et al. 2013,
2015).

The main results can be summarized as follows:

1. Dahlia sits at the centre of a cosmic web knot, and
accretes mass from the intergalactic medium mainly via 3
filaments of length \( \approx 100 \text{kpc} \) (Fig. 4). Dahlia has \( \approx 6 \) major
satellites \( (M_* \approx 10^{10} M_\odot) \) and is surrounded by \( \approx 10 \) minor
ones \((M_* \approx 10^{10} M_\odot)\). The latter represent molecular cloud
(MC) complexes caught in the act of condensing as the gas
streams through the circumgalactic medium (Fig. 5). Dahlia
dominates both the stellar mass \( (M_* \approx 10^{10} M_\odot) \) and the
SFR of the galaxy ensemble \((\text{SFR} \approx 10^2 M_\odot \text{yr}^{-1})\.

2. Only a small fraction of the available energy produced
by stars couples to the gas, as energy is mostly dissipated
within MCs where the stars reside. Radiation dominates the
feedback energy budget by a factor \( > 100 \) (Fig. 8).

3. By \( z = 6 \) Dahlia forms a H2 disk of mass \( M_{\text{H2}} = 3.6 \times
10^{9} M_\odot \), effective radius 0.6 kpc, and scale height 200 pc (Fig.
10). The disk is dense \( (n \approx 25 \text{cm}^{-3}) \), enriched \((Z \approx 0.5 Z_\odot)\),
and it is fed by frequent mergers driving fresh gas to the
centre, and supports a star formation rate per unit area of
\( \approx 15 M_\odot \text{yr}^{-1} \text{kpc}^{-2} \).

4. The disk is mostly unaffected by SN shocks, and it is
pressure-supported by radiation. SN/winds drive hot metal
outflows (Fig. 9), that are either preferentially aligned with
the galaxy rotation axis, or start at the edge of the disk.

5. The total [C II]-luminosity of Dahlia is \( 10^{55.5} L_\odot \), and \( \approx
95\% \) of the emission is co-located with the H2 disk (Fig. 11).
The diffuse, enriched material surrounding Dahlia contains
30\% of the C II mass, but it negligibly contributes to the
[C II] emission (Fig. 10) due to its low density \( (n \approx 10^{-3} \text{cm}^{-3})\)
and metallicity \((Z \approx 10^{-2} Z_\odot)\). Dahlia is under-luminous
with respect to the local [C II]-SFR relation; however, its
luminosity is consistent with upper limits derived for most
\( z \approx 6 \) galaxies.

We find clear indications that the SF subgrid prescrip-
tion might considerably affect the [C II]-SFR relation and
the ISM structure, as noted also by (Hopkins et al. 2013).
This is because stars form in gas of different densities de-
pending on the chosen prescription. In our simulation gas
is converted into stars with an efficiency \( \zeta_{\text{d}} f_{\text{H2}} \), where the
H2 fraction is computed from the KTM09 model and we set
\( \zeta_{\text{d}} = 0.1 \). In S16 the SF follows a total (i.e. not molec-
ular) density Schmidt-Kennicutt relation. Further the SF
efficiency depends on the free-fall time and the turbulent
eddy turnover time. The SF relation is derived from an
empirical fit to MC simulations (Padoan et al. 2012), with no
notion of the local metallicity.

Interestingly, although the approaches are considerably
different, the resulting efficiencies are compatible: in S16 the
bulk of the star forming gas has \( n \approx 10^{1.5} \text{cm}^{-3} \), as in Dahlia
(Fig. 7). However, with respect to S16, Dahlia misses part of
the very dense, star forming gas, and its corresponding
contribution to [C II] from \( Z \approx Z_\odot \) MCs with \( n \approx 10^{2} \text{cm}^{-3} \).
These MC are expected to have high [C II] fluxes (see eq.
We find that the inaccurate computation of the ISM thermodynamic state of a disk would contribute only marginally to the SFR. Thus, unshielded (low dust column density) gas in the H2 clusters might have a strong impact on star formation. For example, Lyman-Werner photons might locally dissociate the H2 by generating pockets of H+ in the gas distribution. This is because the chemical equilibrium assumed in KTM09 does not hold in low-metallicity regimes.

Another important caveat is that we have assumed a uniform UV background. Instead, discrete sources (stellar clusters) might have a strong impact on star formation. For example, very high FUV fluxes can photoevaporate MC on short time scales (≲ tff for gas with Z ~ 10^{-2}Z⊙, Vallini et al. 2016). This effect is particularly important, as it might be responsible for the displacement between the [C ii] and the UV emitting region observed in BDF3299 (Maiolino et al. 2015), and in some of the Capak et al. (2015) galaxies. To solve these problems, a multi-frequency radiative transfer computation must be coupled to the present simulations. This work is ongoing and will be presented elsewhere.

Finally, local FUV flux variations can change the [C ii] emission from individual regions of the galaxy. Also, very high FUV fluxes can photoevaporate MC on short time scales (≲ tff for gas with Z ~ 10^{-2}Z⊙, Vallini et al. 2016). This effect is particularly important, as it might be responsible for the displacement between the [C ii] and the UV emitting region observed in BDF3299 (Maiolino et al. 2015), and in some of the Capak et al. (2015) galaxies. To solve these problems, a multi-frequency radiative transfer computation must be coupled to the present simulations. This work is ongoing and will be presented elsewhere.

ACKNOWLEDGMENTS

We are grateful to the participants of The Cold Universe program held in 2016 at the KITP, UCSB, for discussions during the workshop. We acknowledge the AGORA project members and the DAVID group for stimulating discussion. We thank the authors and the community of PYMES for their work. We thank B. Smith for support in implementing GRACKLE. This research was supported in part by the National Science Foundation under Grant No. NSF PHY11-25915. S.S. was supported by the European Research Council through a Marie-Skodolowska-Curie Fellowship, project PRIMORDIAL-700907.

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**APPENDIX A: BLASTWAVE MODEL**

In this Sec. we present the model used to calculate the kinetic (\( f_{\text{kin}} \)) and thermal (\( f_{\text{th}} \)) energy fractions that a gas cell acquires during a SN explosion/because of a stellar wind (see eqs. 6c).

First, we consider a SN explosion. Then, the picture is the following (e.g. Cioffi et al. 1988; Ostriker & McKee 1988; Walch & Naab 2015). In the first stage of the SN blast the stellar ejecta follow a free expansion solution, that ends when the blast have swept a mass of material roughly the mass of the ejecta, \( \sim 1\text{M}_\odot \) per SN. Then, the shock enters in a Sedov-Taylor stage (ST). The ST stage is adiabatic and the available energy is divided in thermal and kinetic parts, that accounts for \( f_{\text{th}} \simeq 0.7 \) and \( f_{\text{kin}} \simeq 0.3 \) of the total, respectively. The ST stage ends when radiative losses become important, at \( t \approx t_{\text{cool}} \), the cooling time of the ambient gas. At the contact discontinuity the gas cooling causes the formation of a thin shell (shell formation stage, SF), and after that the shock proceeds snowplowing through the ambient medium, driven by the pressure of the gas interior (pressure driven snowplow, PDS). When all the thermal energy is radiated away, the blastwave enters in the so called momentum conserving snowplow (MCS). In the MCS stage the momentum is conserved, while the remaining energy, purely kinetic, is gradually lost because of the work done by the blast on the ambient material, and eventually the blast stops.

In each stage, the blastwave can be modelled by following the analysis by Ostriker & McKee (1988, hereafter OM88). A self similar blastwave can be described by the evolution of the shock front at time \( t \) as \( r_{\text{sh}} \propto t^\eta \), where \( \eta \) is a constant that determines the type of blast (ST, MCS, ...). The velocity of propagation of the shock is \( u_{\text{sh}} = \sigma_r r_{\text{sh}} / t \), and, using the virial theorem (see eq. 3.3 in OM88), the total energy of a blastwave can be written as

\[
E(t) = \left( 4\pi/3 \right) \sigma_r \rho r_{\text{sh}}^2 \propto t^{5/2},
\]

(A1)

where \( \sigma_r \) is a dimensionless constant and \( \rho \) is the density of the material swept by the blast. Within the presented formalism (eq. 6c), \( f = E(t)/E_0 \), where \( E_0 \) is the input energy, i.e. \( E_0 = f_{\text{th}}(t_0, \Delta t) - f_{\text{th}}(t) \) (eq. 5c).

The cooling time is critical in calculating the evolution of the blast between different stages. Here the cooling time \( t_{\text{cool}} \) is defined as the time when \( t = R_0 T_0 / (\pi/2 m_\text{p} \Lambda) \), where \( R_0 \) is the Boltzmann constant, \( T_0 = T_0(t) \) and \( m_\text{p} \) are the temperature and density at the shock front, \( m_\text{p} \) the proton mass and \( \Lambda = m_\text{p} (T_0, Z) \) the cooling function.

Note that calculating the cooling function with GRACKLE or analytical approximations (e.g. Raymond et al. 1976; Nisikawa et al. 1997; Koyama & Inutsuka 2002) yield comparable results. This happens because the gas start to cool at \( T_0 \simeq 10^5 \text{K} \), when Bremsstrahlung is the dominant cooling process, and it is independent of the metallicity, i.e. \( \Delta \propto T^{-1/2} \). Thus, as in Cioffi et al. (1988), see eq 3.10), the cooling time can be approximated as

\[
t_{\text{cool}} \approx 3.6110^{-2} E_0 / [\text{foe}]^{1/4} (n / m_\text{p} \text{cm}^{-3})^{3/4} \text{Myr}
\]

(see also Kim & Ostriker 2015, in particular sec. 6 and 7).

The typical range of input energy (\( 1 \lesssim E_0 [\text{foe}] \lesssim 10^3 \)) and ISM density (\( 10^{-2} \lesssim \rho / m_\text{p} \text{cm}^{-3} \lesssim 10^2 \)), thus the cooling time is in the range \( 10^{-3} \lesssim t_{\text{cool}} \text{Myr} \lesssim 1 \). Since the expected simulation time step is \( \Delta t \sim 10^{-2} \text{Myr} \), we can further simplify the blastwave picture.

The free expansion stage is shorter than our typical simulation time step (\( t \lesssim 10^{-4} \text{Myr} \), e.g. see eq. 1 in Kim & Ostriker 2015), thus we assume that the SN starts in the ST stage (\( \eta = 2/5 \)). After \( t \approx t_{\text{cool}} \), the energy of the shock is roughly half the initial and the blastwave is in the PDS stage (\( \eta = 2/7 \)). In both stages, the total energy is given by the sum of kinetic and thermal terms, and the relative fraction of kinetic (\( f_{\text{kin}} \)) and thermal (\( f_{\text{th}} \)) energies are constants (see eqs. 3.16 and 3.18 in OM88) for both. In the intermediate SF stage we approximate \( f_{\text{kin}} \) and \( f_{\text{th}} \) by linearly interpolating between the ST and initial PDS values. The time when the blast enters in the SF and PDS stages can be calculated following the analytical fit to the simulation presented in Cioffi et al. (1988), i.e. \( 0.14 t_{\text{cool}} \) and \( 0.4 t_{\text{cool}} \) for SF and PDS, respectively (see their eq. 3.15). An example of such model is presented in Sec. 2.4 (in particular, see Fig. 3). Note that the model is consistent with the result from SN exploding in an homogeneous medium, however a blastwave propagating in an inhomogeneous medium behave differently, since the blast travels unimpeded through path of lower density medium (Martizzi at al. 2015, see in particular the fit in eq. 9 and 10 and the different scalings given in eq. 11 and 12). We will explore this effect in a future work.

The winds can be treated within the same blastwave formalism (see Sec. VII in OM88). The stage evolution of the winds is similar to the SN one: first the wind is adiabatic, then the outer shock begin to radiate away and a thin shell is formed, for the complete picture of blastwave evolution, we refer the reader to OM88, in particular Fig.1, table IV and references therein.
and eventually the shock becomes momentum conserving. However in the wind case the energy injection is not impulsive, as the stars input a continuous energy injection with a (roughly) constant luminosity. Thus, we have a different blastwave structure and consequently the stant luminosity. Thus, we have a different blastwave structure and eventually the shock becomes momentum conserving. How-

ever in the wind case the energy injection is not impulsive, as the wind solution is given by Weaver et al. (1977, hereafter W77). The structure can be spatially divided in three parts (see Fig. 1 in W77): near to the stars we have the stellar wind (α), and the gas is contained in the region of shocked stellar wind (b) and in the shell of shocked gas (c).

In the adiabatic case we have a (roughly) constant wind luminosity (L) and no radiative losses, thus dimensional analysis implies η = 3/5 in eq. A1. In this stage the gas in region 6 has 5/11 of the total energy, that is in a purely thermal form, while the gas in c contains energy in both thermal and kinetic form. Summing the contribution of both regions, W77 finds f_b ≈ 0.78 f_kn ≈ 0.22 (see also eqs. 7.11 in OM88). Note that the thermal energy fraction is larger to the corresponding one in the ST stage of the SN driven blast.

In the next stage radiative losses becomes important (t > t_coal), and region (c) collapses to a thin shell. All the energy (5/11 Lt) would be contained in the gas in b in thermal. The physical situation is that the injected fluid is adiabatic, while in the ambient radiative losses would be dominant. This happens because the density of the injected fluid is much less then the ambient medium, and this consequently affects the the cooling time scales. However, the modellization of this stage cannot be readily implemented as a subgrid model in our simulation, because we cannot easily keep track of the internal structure of the cells in the simulation.

In our code we opt to go directly from the adiabatic stage (f_b = 0.78, f_kn = 0.22, W77) to the momentum conserving stage when t > t_coal. The latter is described by f_kn = 1, and E(t) is given by eq. A1 with η = 1/4 (see eq. 7.20 in OM88).

As noted in Agertz et al. (2013); Walsh & Naab (2015); Fierlinger et al. (2016); Körtgen et al. (2016), we expect wind injection to make the SN more efficient, since SN blast sweeps through a gas with a lower density, thus with a longer cooling time. This point applies both to wind and radiation pressure injection.

Finally, note that in the blastwave modellization, we have neglected the instabilities that can arise during the thin shell formation stage (Madau et al. 2001; McLeod & Whitworth 2013; Badjin et al. 2016), and we do not explicitly consider the effect of multiple blastwave events (Walsh & Naab 2015; Fierlinger et al. 2016). This issues will be addressed by future work.

APPENDIX B: RADIATION PRESSURE: OPTICAL DEPTH AND IR-TRAPPING

The radiation pressure is dependent on the rate of momentum injection (p_rad, eq. 7a), that in turn depend on the optical depth to ionizing photons (τ_{ion}) and the IR-trapping of the UV radiation (f_{ir}).

For each gas cell, the optical depth τ_{ion} can be calculated by averaging the ionization cross section on the stellar spectra assumed in our model (see eqs. 5e), that are taken from STARRHURST99 (Leitherer et al. 1999, 2010). Namely, τ_{ion} = σ_{ion}N_H, where ionization cross-section σ_{ion} is obtained as a Rosseland mean

\[ \sigma_{ion} = \int_{1000}^{912} \frac{\rho}{L} \lambda^2 \sigma_{ion} \, d\lambda / L_{ion}. \]

where the cross section as a function of wavelength is given by Osterbrock (1989)

\[ \sigma_{ion} = 6.3 \times 10^{-18} (1.34 (912/\lambda)^{-2.99} + 0.34 (912/\lambda)^{-3.99}). \]

For the IR-trapping we assume f_{ir} = τ_{ir}, the infrared dust optical depth, similarly to Agertz et al. (2013). As noted in Krumholz & Thompson (2012), the assumption f_{ir} = τ_{ir} is approximately correct for small values of the optical depth, i.e. τ_{ir} ≲ 20. This is our case, since in simulations with resolution scale larger than 10pc, we expect τ_{ir} ≲ 10 (Rosdahl et al. 2015). As reported in Sec. 3.2.2, τ_{ir} ≲ 10^{-1} in our simulation (see also Fig. 8).

Using the cross section from Draine (2003, see also Semenov et al. 2003), we can write

\[ \tau_{ir} \approx 7.8 \times 10^{-4} (L_H / L_{\odot})^{1/6} (M_{dust} / M_{\odot})^{-1/6}, \]

(B2a)

where \( T_{dust} \) is the dust temperature. \( T_{dust} \) can be calculated as (Dayal et al. 2010)

\[ T_{dust} = 6.73 (L_H / L_{\odot})^{1/6} (M_{dust} / M_{\odot})^{-1/6}, \]

(B2b)

where \( L_H \) is the reprocessed UV luminosity, i.e. \( L_H = L_{uv}(1 - \exp(\tau_{uv})) \), and \( M_{dust} \) is the dust mass, i.e. \( M_{dust} = \rho_{dust} \nu_{dust} (2 \pi Z / X_{dust}), \)

(B2c)

with \( \rho_{dust} = 6 \times 10^{-3} \) being the solar dust to gas ratio (e.g. Hirashita & Ferrara 2002).

We account for CMB heating on dust, that can be important at high redshift (e.g. da Cunha et al. 2013). The CMB heating is calculated by using a correction to the dust temperature (da Cunha et al. 2013, in particular see eq. 12), i.e.

\[ T_{dust}^{corr}(z) = (T_{dust}^{4+\beta} + \frac{4+\beta}{4+\beta} (\tau_{CMB} - T_{dust}^{4+\beta} z = 0))^{1/(4+\beta)} \]

(B2d)

where \( T_{CMB}(z) = 2.725 (1 + z) \) is the average CMB temperature at \( z \) and \( \beta \) is the dust emissivity coefficient (Draine & Lee 1984). Similarly to da Cunha et al. (2013), we assume a fiducial dust emissivity of \( \beta = 2 \).

Note that dust can be destroyed by sublimation or evaporation. In the adopted chemical network (and the current CRACKLE version), dust abundance is not accounted self consistently, and we calculate it by assuming that the \( M_{dust} \) is proportional to the metal mass of the gas (eq. B2c). To mimic dust destruction, we assume a negligible dust contribution to the optical depth when \( T_{dust} > 2 \times 10^{4} \) (e.g. Bauer et al. 1997). We neglect dust sputtering driven by SNe (e.g. Valiante et al. 2009; Draine 2011), which might lead to dust destruction, and thus further reduce the radiation pressure efficiency.

Finally, we note that in our radiation feedback model (eqs 7a and 7c), energy conservation is guaranteed by construction, if only the UV and ionizing contributions were present. However, the IR-trapping is accounted by approximate formulas (f_{ir} = τ_{ir}), that are consistently evaluated with the gas properties but do not account for radiative losses in the IR cascade. For this reason, we have added the additional energy conservation check presented in Sec. 2.4.3. Note that a posteriori we have found that IR-trapping is subdominant contribution to radiation pressure (τ_{ir} ≲ 10^{-1}, Sec. 3.2.2), thus probably the check is not necessary.

APPENDIX C: POSTPROCESSING IONIZATION STATE AND EMISSION

Radiative transfer is not followed during the evolution of the simulation. To compute the ionization state of the various atomic species, we post-process the simulation outputs using the photoionization code CLOUDY (Ferland et al. 2013). We consider a grid of models based on the density (n), temperature (T) and metallicity (Z) of the gas in our simulation. We produce a total of 10^9 models, that are parameterized as a function of the column density (N).

The radiation fields includes the UVB intensity at 912 Å (Haardt & Madau 2012), the CMB background and a galactic background, that is obtained by rescaling with Dahlia SFR the Galaxy spectrum (Black 1987), in particular G the FUV flux in the Habing band, that is usually normalized to the Galactic mean.
value $G_0$. Thus we use a use $G = 130 G_0$ in our calculation. Note that larger value of $G$ does not yield a large variation of the expected [C II] in molecular gas (Vallini et al. 2016).

Note that accounting for the UVB is not relevant for the ionization state of the gas in the proximity of galaxies (Gnedin 2010): because of the high density environment, the gas is efficiently shielded and the galactic emission is the dominant radiation source. Thus, in our cloudy models we consider that the gas is shielded by a column density of $N \simeq 10^{20}$ cm$^{-2}$.

Spatial variation of the incident radiation are are not considered in the present work. It is to note that the variation of $G$ is very low in our Galaxy (Habing 1968; Wolfire et al. 2003), i.e. $(G) = G_0$ and $(\langle (G/G_0)^2 - 1 \rangle)^{1/2} \simeq 3$. However, in the close proximity of young OB associations, the flux can be very high, e.g. $10^6 G_0$ at 0.1 pc from a single OB star (Hollenbach & Tielens 1999), or – equivalently – $10^6 G_0$ at 10 pc from a starburst with $SFR \simeq M_\odot$/yr. As the flux propagate with $1/r^2$, such variation are not perceived in most of the volume. Such effect is not considered in the present work.

Using the C II density obtained with cloudy, we compute the [C II] luminosity by using eq. 3 in Vallini et al. (2013). The effect of CMB suppression of [C II] is included in the model (Pallottini et al. 2015; Vallini et al. 2015). Such effect suppress the emission where the spin temperature of the [C II] transition is close to the CMB one. This is relevant for low density ($n < \sim 10^{-3}$ cm$^{-3}$) medium, that does not have enough collision to decouple from the CMB. We note that the result from the present [C II] modelling is consistent with the one obtained by directly using cloudy to compute the PDR emission.

Note that the current model does not account for the photoevaporation effect on MC, that has an important impact on FIR emission (Vallini et al. 2016), particularly when including a spatially varying FUV field. More detailed modelling will be accounted in a future work.