

New composition-dependent cooling and heating curves for galaxy evolution simulations

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ABSTRACT

In this paper, we present a new calculation of composition-dependent radiative cooling and heating curves of low-density gas, intended primarily for use in numerical simulations of galaxy formation and evolution. These curves depend on only five parameters: temperature, density, redshift, [Fe/H] and [Mg/Fe]. They are easily tabulated and can be efficiently interpolated during a simulation. The ionization equilibrium of 14 key elements is determined for temperatures between 10 K and 10^9 K and densities up to 100 amu cm^{-3} taking into account collisional and radiative ionization, by the cosmic UV background and an interstellar radiation field, and by charge-transfer reactions. These elements, ranging from H to Ni, are the ones most abundantly produced and/or released by SNIa, SNII and intermediate-mass stars. Self-shielding of the gas at high densities by neutral hydrogen is taken into account in an approximate way by exponentially suppressing the H-ionizing part of the cosmic UV background for H I densities above a threshold density of $n_{\text{H I, crit}} = 0.007 \text{ cm}^{-3}$. We discuss how the ionization equilibrium, and the cooling and heating curves, depends on the physical properties of the gas. The main advantage of the work presented here is that, within the confines of a well-defined chemical evolution model and adopting the ionization equilibrium approximation, it provides accurate cooling and heating curves for a wide range of physical and chemical gas properties, including the effects of self-shielding. The latter is key to resolving the formation of cold, neutral, high-density clouds suitable for star formation in galaxy simulations.

Key words: atomic processes – hydrodynamics – plasmas – ISM: general.

1 INTRODUCTION

Numerical simulations of galaxy evolution require basic physical input regarding the (thermo-)dynamical behaviour of the interstellar gas. A crucial ingredient of the energy, or entropy, equation is the cooling rate of the gas. This quantity is, in principle, a complex function of the temperature, composition and irradiation of the gas. An often used assumption is that the gas is in collisional ionization equilibrium (CIE). In that case, collisions with free electrons are deemed solely responsible for keeping atoms ionized. Since both the recombination rate and the ionization rate are in that case directly proportional to the electron density, the latter cancels from the equations and the ionization equilibrium becomes a function of temperature only (for a given elemental abundance mix). For low gas densities, each collisional ionization/excitation is followed by a

radiative de-excitation, creating an escaping photon, and the cooling rate becomes proportional to the density squared (or to hydrogen density times electron density, $n_{\text{H}}n_{\text{e}}$) times a temperature-dependent function.

Many state-of-the-art simulation codes (Revaz et al. 2009; Sawala et al. 2011; Scannapieco et al. 2011; Schroyen et al. 2011; Cloet-Osselaer et al. 2012; Gabor & Davé 2012; Kim et al. 2012) rely on the cooling curves compiled by Sutherland & Dopita (1993). The latter authors calculated cooling rates, excluding a forefactor $n_{\text{H}}n_{\text{e}}$, as a function of temperature for a number of metallicities. During a simulation, the cooling rate of a gas parcel can be rapidly determined by simple two-dimensional interpolation on these curves. However, while this work was monumental and has spawned a large volume of literature based on simulations making use of these curves, one needs to be aware of the assumptions on which these cooling curves are based and simulators need to assess whether they can be used for the application at hand. To be clear: this is no criticism of the Sutherland & Dopita (1993) cooling curves.

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(i) Metallicity is quantified using $[\text{Fe}/\text{H}]$, the iron abundance. For $[\text{Fe}/\text{H}] \leq -1$ the abundance ratios are taken to reflect those of a SNII, with $[\alpha/\text{Fe}] > 0$; for $[\text{Fe}/\text{H}] = 0$ the solar abundance ratios are adopted, with $[\alpha/\text{Fe}] = 0$; for other metallicities, the abundance ratios are interpolated linearly between these two sets of abundance ratios. Hence, *adopting the Sutherland & Dopita (1993) cooling curves immediately implies adopting the solar neighbourhood's chemical enrichment history*. In some cases, this may not be a good approximation of reality. For example, dwarf galaxies have low metallicities but also low $[\alpha/\text{Fe}]$ (see e.g. Tolstoy, Hill & Tosi 2009, and references therein). Using the low metallicity (Sutherland & Dopita 1993) cooling curves will then quite strongly overestimate the cooling contributed by α -elements such as O, Si and Mg. The centres of giant elliptical galaxies, on the other hand, have high metallicities and high $[\alpha/\text{Fe}]$ (Worthey, Faber & Gonzalez 1992). With the Sutherland & Dopita (1993) cooling curves, the contribution of the α elements will be strongly underestimated.

(ii) Another issue is whether in the presence of a cosmological UV background (UVB) CIE is still an acceptable approximation. The UVB tends to keep (part of) the hydrogen and metals ionized, thus lowering the fraction of H I and raising n_e . This dramatically influences the shape of the cooling curve, as is well known (Wiersma, Schaye & Smith 2009) and as we will also show below. Tepper-García et al. (2011) compared the net cooling rates computed assuming only CIE with those calculated including photo-ionization. These authors showed that the equilibrium temperature of the gas could be off by an order of magnitude at low densities and high metallicities when using CIE.

(iii) Simulations nowadays achieve sufficiently high spatial resolution to be able to follow the formation of cold, dense star-forming clouds. This requires extending the Sutherland & Dopita (1993) cooling curves to temperatures below 10^4 K. An often used extension is the set of cooling curves of Maio et al. (2007). These authors calculate the level populations of Fe II , O I , Si II and C II and the cooling rates due to the low-lying finestructure emission lines of these ions. These level populations are set by collisions with free electrons and thus by the ionization fraction of the gas. Again, this is a quantity which is very sensitive to the presence of a cosmic UVB. Clearly, the often used approach of ‘gluing’ the Maio et al. (2007) cooling curves ($10 < T < 10^4$ K) to the Sutherland & Dopita (1993) cooling curves ($10^4 < T < 10^9$ K) will produce unreliable results in simulations with a cosmic UVB.

Recent calculations of cooling curves with `CLOUDY`, such as Wiersma et al. (2009) (used in the Overwhelmingly Large Simulations project; Schaye et al. 2010) and Shen, Wadsley & Stinson (2010) [used in the simulation code `GASOLINE` (Shen et al. 2010; Brook et al. 2012)], include the cosmic UVB in determining the ionization balance. Here, the total cooling rate is written as the sum of approximately independent terms: the cooling due to H and He, the cooling due to metals and inverse Compton cooling. Wiersma et al. (2009) advocate an element-by-element approach, necessitating the tracing of the abundances of a set of 11 elements during a simulation in order to calculate the independent contribution of each to the total cooling rate. In reality, these terms are linked by the free electron density, by charge-exchange reactions, and by other reactions between different elements (such as molecule and dust particle formation). Moreover, many authors still adopt the solar abundance ratios and scale the metal cooling rate proportional to metallicity which, as we have argued, can be expected to be a bad representation of reality for certain types of stellar systems. Likewise, Smith, Sigurdsson & Abel (2008), and Smith et al. (2011), only provide

cooling curves for solar abundance ratios of the heavy elements (H and He ions are followed explicitly during simulations). In the former paper, the influence of the UVB is not taken into account while the latter focuses precisely on this issue. Gnedin & Hollon (2012) incorporate the radiation field through four well-chosen normalized photoionization rates. This, together with a Taylor expansion of the curves up to quadratic terms in metallicity, yields an approximation to the cooling and heating curves with a median error of 10 per cent but with (although very rare) errors of up to a factor of 6.

In this paper, we try to improve on several aspects of the existing cooling curve calculations.

2 NUMERICAL DETAILS

Below, we give a list of the most prominent ingredients of our calculations:

(i) We adopt a chemical enrichment model that is self-consistent in the sense that, in an N -body/SPH simulation, gas particles can be enriched by stellar particles in only two ways: fast [by SNII and massive intermediate-mass stars (IMS), with $M \sim 8 M_\odot$] and slow (by SNIa and less massive IMS). Thus, the chemical abundance mix of a gas particle depends solely on the ratio of the ‘slow’ and ‘fast’ contributions. The cooling and heating rates can then be tabulated for a small number of different ratios of ‘slow’ to ‘fast’ contributions, covering all possibilities that can occur in a simulation.

SNII yields are taken from Nomoto et al. (1997) and modified according to the prescriptions detailed in François et al. (2004); IMS yields come from Gavilán, Buell & Mollá (2005); SNIa yields have been adopted from Tsujimoto et al. (1995). For SNII, the fraction of the initial mass of a stellar population that is returned in the form of element X to the interstellar medium (ISM) is given by

$$y_{X,\text{SNII}} = \frac{\int_{m_{\text{III}}}^{m_{\text{uII}}} M_X(m)\phi(m)dm}{\int_{m_1}^{m_u} m\phi(m)dm}. \quad (1)$$

Here, $m_{\text{III}} = 8 M_\odot$ and $m_{\text{uII}} = 70 M_\odot$ are the lower and upper bounds of the masses of stars that turn into SNII; $m_1 = 0.1 M_\odot$ and $m_u = 70 M_\odot$ are the adopted lower and upper bounds of the masses of stars. The mass returned in the form of element X by a star with initial mass m is denoted by $M_X(m)$. For the initial-mass function, or IMF, denoted by $\phi(m)$, we take the parametrization by Chabrier (2003).

For SNIa, the fraction of the initial mass of a stellar population that is returned in the form of element X to the ISM is given by

$$y_{X,\text{SNIa}} = A_{\text{Ia}} M_X \frac{\int_{m_{\text{Ia}}}^{m_{\text{III}}} \phi(m)dm}{\int_{m_1}^{m_u} m\phi(m)dm}. \quad (2)$$

Here, $m_{\text{Ia}} = 3 M_\odot$ is the lower bound of the masses of stars that can turn into SNIa and M_X is the SNIa yield of element X . The forefactor A_{Ia} was determined by demanding that the calculated ratio of the occurrence of SNIa to that of SNII reproduces that derived for the solar neighbourhood by Tsujimoto et al. (1995):

$$\frac{N_{\text{Ia}}}{N_{\text{II}}} = 0.15 = A_{\text{Ia}} \frac{\int_{m_{\text{Ia}}}^{m_{\text{III}}} \phi(m)dm}{\int_{m_{\text{III}}}^{m_{\text{uII}}} \phi(m)dm}. \quad (3)$$

For IMS, the yield is given analogously by

$$y_{X,\text{IMS}} = \frac{\int_{m_{\text{IMS}}}^{m_{\text{uIMS}}} M_X(m)\phi(m)dm}{\int_{m_1}^{m_u} m\phi(m)dm}, \quad (4)$$

Table 1. Element yields.

Element	Slow	Fast
O	0.000136	0.000937
C	0.000146	0.000143
Ne	2.41e-05	0.00013
Mg	1.06e-05	6.23e-05
Si	1.66e-05	4.89e-05
Fe	2.96e-05	1.67e-05
S	9.15e-06	1.41e-05
N	5.87e-05	9.65e-06
Al	6.46e-08	6.51e-06
Na	4.86e-09	3.36e-06
Ni	2.99e-07	1.67e-06
Ca	9.99e-07	1.64e-06

with $m_{\text{IMS}} = 0.8 M_{\odot}$ and $m_{\text{ulIMS}} = 8 M_{\odot}$. The yields of elements contributed by the most massive IMS, such as ^{13}C and N, are added to the corresponding SNII yields. Those of elements produced by longer-lived stars are added to the corresponding SNIa yields. This way, there are two contributions to the yield of a given element: a ‘fast’ one (encompassing the contributions from SNII and massive IMS) and a ‘slow’ one (encompassing the contributions from SNIa and less massive IMS). The abundance of each chemical element in a gas parcel is then simply the weighted sum of these two contributions.

Since this simple chemical evolution model contains two sets of yields, the elemental abundance ratios in a given gas parcel can also be quantified by just two numbers. Here, we choose $[\text{Fe}/\text{H}]$ as a tracer of the overall metallicity and $[\text{Mg}/\text{Fe}]$ as a second parameter. Mg is an α -element released abundantly by SNII explosions but produced in only very small amounts by SNIa explosions so it is a good tracer of the relative weights of the ‘fast’ and ‘slow’ contributions. Moreover, there are now quite advanced techniques available to determine the abundances of both Fe and Mg in a stellar population from absorption lines in optical spectra. Obviously, oxygen would also make a good discriminator between the ‘fast’ and ‘slow’ contributions. We present the element yields used in this paper, expressed as a fraction of the initial mass of a single stellar population, in Table 1.

In Fig. 1, we compare the observed abundances of a number of well-studied stars, taken from Worley et al. (2009), Grevesse et al. (2010), Fitzpatrick (2010), Landstreet (2011), with those predicted by our chemical evolution model for the appropriate $[\text{Fe}/\text{H}]$ and $[\text{Mg}/\text{Fe}]$ values. For the Sun and Arcturus, the largest discrepancies are about 0.2 dex. For Sirius and Vega, there are larger deviations between data and model, although the author-to-author scatter on the measured abundances of these stars is, admittedly, quite substantive (e.g. the Na abundance in the atmosphere of Sirius varies by more than one dex between different authors; see e.g. Landstreet 2011). Taking into account observational uncertainties and genuine cosmic scatter, this is reassuring evidence that the simple two-yield chemical evolution model employed here works adequately.

(ii) The ionization balance, electron density, level populations and cooling rates are calculated self-consistently in the presence of a cosmic UVB. Here, we adopt the UVB calculated by Faucher-Giguère et al. (2009). We use the UV spectra available from this author’s webpage¹ to calculate the ionization and heating rates of all elements.

(iii) Stars also generate an interstellar radiation field (ISRF), capable of ionizing atoms with small ionizing potentials, such as C I, Si I, Mg I, Ca I, Ca II, Fe I, etc. even at very low gas temperatures. Since the light of newly formed stars had to make its way through the H II regions surrounding these stars, it no longer has an H-ionizing component. We include the parametrized ISRF of Mathis, Mezger & Panagia (1983), appropriate for the solar neighbourhood. The ISRF’s main task is to keep the elements mentioned above ionized while its precise form has shown to be of little consequence. At low gas densities, the UVB is the dominant photo-ionizing radiation field while at high densities, where star formation becomes important and neutral hydrogen can shield the gas from the UVB, the ISRF gains importance.

(iv) Through hydrogen ionizations and cascade recombinations, the H-ionizing portion of diffuse UV radiation impacting on a gas cloud is converted into lower-frequency radiation. Thus, for sufficiently high densities, gas may become self-shielding against H-ionizing UV radiation once hydrogen recombinates. Therefore, self-shielding will generally be insignificant for temperatures $T \gtrsim 10^4$ K since then hydrogen is collisionally ionized anyway.

This is not a straightforward problem since it in principle requires solving the radiative transfer equation. However, one can estimate the critical H density above which self-shielding can be expected to block most of the ionizing UV radiation. Tajiri & Umemura (1998) put forward $n_{\text{H}} \sim 0.01 \text{ cm}^{-3}$, Aubert & Teyssier (2010) estimate that $n_{\text{H}} = 0.007 \text{ cm}^{-3}$, and Yajima, Jun-Hwan & Nagamine (2011) quote $n_{\text{H}} = 0.00634 \text{ cm}^{-3}$. With this cut-off density, it is possible to reproduce the observed mass- and volume-averaged neutral fraction of the Universe at a redshift $z \sim 6$ (Aubert & Teyssier 2010), the H I column density distribution of damped Ly α systems $z = 3$ (Nagamine, Choi & Yajima 2010) and the Ly α luminosity of forming galaxies (Faucher-Giguère et al. 2010).

We have implemented an approximate scheme for self-shielding by exponentially suppressing the UV radiation field with frequencies above $h\nu = \chi_{\text{H I}}$, with $\chi_{\text{H I}}$ the hydrogen ionization potential, as

$$J_{\nu}(\nu, n_{\text{H I}}) = J_{\nu}(\nu) \exp(-n_{\text{H I}}/n_{\text{H I, crit}}) \quad h\nu > 1 \text{ Ry} \\ = J_{\nu}(\nu) \quad h\nu \leq 1 \text{ Ry}, \quad (5)$$

with $n_{\text{H I, crit}} = 0.007 \text{ cm}^{-3}$ and $J_{\nu}(\nu)$ the original UV spectrum, as in Aubert & Teyssier (2010). Note that we use the *neutral* H density here, not the *total* H density, since it is only the neutral fraction which is responsible for absorbing H-ionizing UV radiation.

One could worry that, when the gas at high densities and low temperatures becomes self-shielding against the external UVB, the cooling radiation itself may become trapped and be re-absorbed, affecting the cooling rate and the ionization equilibrium (Gnat & Sternberg 2007). However, in a self-shielding H I cloud below $T \sim 10^4$ K, only low-energy UV photons unable to photo-heat the hydrogen gas are emitted. Moreover, with most of the star formation prescriptions currently popular in galaxy evolution and cosmological simulations (Governato et al. 2010; Schroyen et al. 2011; Cloet-Osselaer et al. 2012), such clouds will begin to form stars before reaching densities exceeding 100 amu cm^{-3} and stellar and supernova feedback will rapidly overwhelm any internal diffuse radiation field. Therefore, we expect this to be a minor issue.

(v) Charge-exchange reactions can efficiently transfer electrons between ions with similar ionization potentials. Given their high abundances, H I and H II are the ions’ most likely reaction partners. Some of the ions that play an important role in gas cooling below 10^4 K via fine-structure line emission, such as O I, are

¹ <https://www.cfa.harvard.edu/~cgiguere/UVB.html>

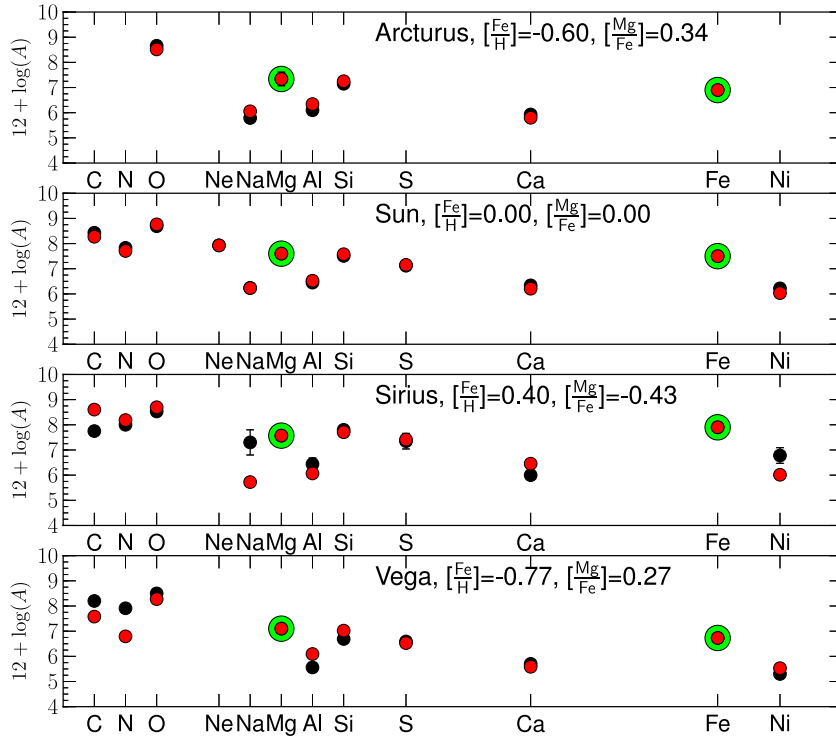


Figure 1. The observed metal abundances in the atmospheres of a number of well-studied standard stars (black points) and those predicted by our chemical evolution model (red points) for the appropriate $[\text{Fe}/\text{H}]$ and $[\text{Mg}/\text{Fe}]$ values (green points). The abundance, A , of an element is measured as the ratio of its number density relative to that of H.

particularly affected by these reactions. We adopt the charge-transfer reaction rates for C I, C II, O I–V, Si I–V and Fe I–V from Kingdon & Ferland (1996) and Stancil et al. (1998, 1999) and from the online ORNL/UGA Charge Transfer Database for Astrophysics (<http://www-cfadc.phy.ornl.gov/astro/ps/data/>).

(vi) Charge-exchange reactions are but one example of reactions that involve ions of different elements. Other examples are molecule and dust particle formation. We opted not to include molecular processes in the present work for the following reasons. Judging from e.g. Maio et al. (2007) and Vasiliev (2013), cooling below $T \sim 10^3$ K by molecules is dominated by cooling by metals once the latter are present at levels $Z \gtrsim 10^{-3} Z_{\odot}$. However, at the low metallicities where H_2 cooling might be important, the low dust content strongly inhibits the formation of H_2 and the time-scale for the conversion of H I to H_2 becomes much larger than the local free-fall time (Krumholz 2012). Hence, in this metallicity range, star formation will precede the formation of H_2 . Also, the rates at which dust forms and at which H_2 molecules form via grain catalysis contain many extra free parameters (such as the time-scales for dust formation, growth, and destruction, the dust grain size, the probability for an H atom to stick to a grain, the probability that two H atoms on a grain join and detach themselves from the grain as a single H_2 molecule, etc.) that need to be constrained by experiments and observations (Bekki 2013). Still, by neglecting molecular cooling we might be under-estimating the cooling rate at temperatures below $T \sim 10^3$ K for metallicities below $Z \sim 10^{-3} Z_{\odot}$. However, by not using the ‘independent element’ approximation, as in Wiersma et al. (2009) and Shen et al. (2010), to calculate the ionization balance and the cooling curves, our approach can straightforwardly be extended in future work.

(vii) Using all these ingredients, we calculate cooling and heating curves for the temperature range $10 \text{ K} < T < 10^9 \text{ K}$. This way,

the cooling and heating rates are calculated in perfect consistence with the ionization equilibrium over a very wide range of temperatures. Thus, there is no need to stitch together cooling curves from different authors with potentially very different (and inconsistent) input physics.

For this work, we have extended the capabilities of CHIANTIPY, a PYTHON interface to the CHIANTI atomic data base (Dere et al. 2009), available from <http://chiantipy.sourceforge.net/>. For all ions, we use the recombination rates, collisional ionization rates and energy level populations provided by standard CHIANTIPY. Photo-ionization cross-sections $\sigma_i(\nu)$ are adopted from Verner et al. (1996) and integrated over the stellar and cosmic UVBs in order to obtain the photo-ionization rate

$$\Gamma_i = 4\pi \int_{\nu_i}^{\infty} \sigma_i(\nu) J_{\nu}(\nu) \frac{d\nu}{h\nu}, \quad (6)$$

and the photo-heating rate

$$\dot{q}_i = 4\pi \int_{\nu_i}^{\infty} \sigma_i(\nu) J_{\nu}(\nu) (h\nu - h\nu_i) \frac{d\nu}{h\nu}, \quad (7)$$

with ν_i the ion’s ionization threshold. The integral over the radiation backgrounds is split in two parts: the part for photon energies above 1 Ry, which can be suppressed by H I self-shielding, and the part for photon energies below 1 Ry, which is assumed to be unaffected by self-shielding. For instance, the photo-ionization rate of an element can then be written as

$$\Gamma_i = e^{-n_{\text{H1}}/n_{\text{H1,crit}}} 4\pi \int_{1\text{Ry}}^{\infty} \sigma_i(\nu) J_{\nu}(\nu) \frac{d\nu}{h\nu} + 4\pi \int_{\nu_i}^{1\text{Ry}} \sigma_i(\nu) J_{\nu}(\nu) \frac{d\nu}{h\nu}. \quad (8)$$

For a given temperature, the ionization equilibrium, i.e. the density of the r -times ionized ion of any element with atomic number Z , denoted by $n_{Z,r}$, is found employing a multi-dimensional Newton–Raphson technique using back-tracking. At each iteration, the electron density is given by

$$n_e = \sum_{Z \geq 1} \sum_{r=0}^Z r n_{Z,r}. \quad (9)$$

The general equilibrium condition then becomes

$$\begin{aligned} n_{Z,r} \left[n_e \text{Rec}_{Z,r}(T) + \sum_c n_c \text{lon}_{Z,r}^c(T) + \Gamma_{Z,r} \right. \\ \left. + n_{\text{H I}} \text{CT}_{Z,r}^{\text{H I}}(T) + n_{\text{H II}} \text{CT}_{Z,r}^{\text{H II}}(T) \right] \\ = n_{Z,r+1} \left[n_e \text{Rec}_{Z,r+1}(T) + n_{\text{H I}} \text{CT}_{Z,r+1}^{\text{H I}}(T) \right] \\ + n_{Z,r-1} \left[\sum_c n_c \text{lon}_{Z,r-1}^c(T) + n_{\text{H II}} \text{CT}_{Z,r-1}^{\text{H II}}(T) + \Gamma_{Z,r-1} \right]. \end{aligned} \quad (10)$$

Here, Rec indicates the ionic recombination rate, lon^c represents the collisional ionization rate with collisional partner c (which could be electrons, protons, hydrogen atoms, . . .), and CT stand for charge-transfer reaction rates (some of which are obviously zero, such as the reaction rate between H I and a neutral atom). Given the strong, in this case exponential, dependence of the self-shielding on the neutral hydrogen fraction this is clearly a very non-linear set of equations.

Moreover, the UVB, at least at sufficiently low gas densities, can be expected to keep a large fraction of the hydrogen gas ionized, thereby suppressing the bound–bound and free–bound cooling contributed by hydrogen. As a consequence, the pronounced peak in the CIE cooling rate around $T \sim 10^4$ K, which is caused by hydrogen, will be absent. Since non-equilibrium cooling generally leads to over-ionization compared with CIE, it also tends to suppress the H cooling peak. Hence, the ionization equilibrium assumption will have a smaller impact on cooling rates calculated in the presence of a UVB than on cooling rates calculated assuming CIE (Wiersma et al. 2009). As a check on the assumption of ionization equilibrium, we calculated the ionization time-scales, τ_{ion} , of e.g. the carbon ions to be compared with the cooling time-scale, τ_{cool} . The former is given by

$$\begin{aligned} \frac{1}{\tau_{\text{ion}}} \approx n_e \text{Rec}_{Z,r}(T) + \sum_c n_c \text{lon}_{Z,r}^c(T) + \Gamma_{Z,r} \\ + n_{\text{H I}} \text{CT}_{Z,r}^{\text{H I}}(T) + n_{\text{H II}} \text{CT}_{Z,r}^{\text{H II}}(T) \end{aligned} \quad (11)$$

while for the latter we use

$$\tau_{\text{cool}} = \frac{3}{2} \frac{nkT}{|\Lambda(T) - \mathcal{H}(T)|} = \frac{3}{2} \frac{nkT}{\Lambda_{\text{net}}(T)} \quad (12)$$

with $\Lambda(T)$ the cooling rate and $\mathcal{H}(T)$ the heating rate. The ionization equilibrium approximation is valid if $\tau_{\text{cool}} \gg \tau_{\text{ion}}$ for all ions. Judging from Fig. 2, this constraint is more easily fulfilled at low densities, when the UVB irradiates the gas unimpeded and keeps most of the hydrogen ionized. Hence, the net cooling rate $\Lambda_{\text{net}}(T)$ is small. At higher densities the UVB is attenuated, hydrogen recombines and $\Lambda_{\text{net}}(T)$ is large. This makes it much harder for this constraint to be fulfilled, especially at lower temperatures. This is a caveat that should, of course, be kept in mind when using any set of cooling and heating tables calculated assuming ionization equilibrium.

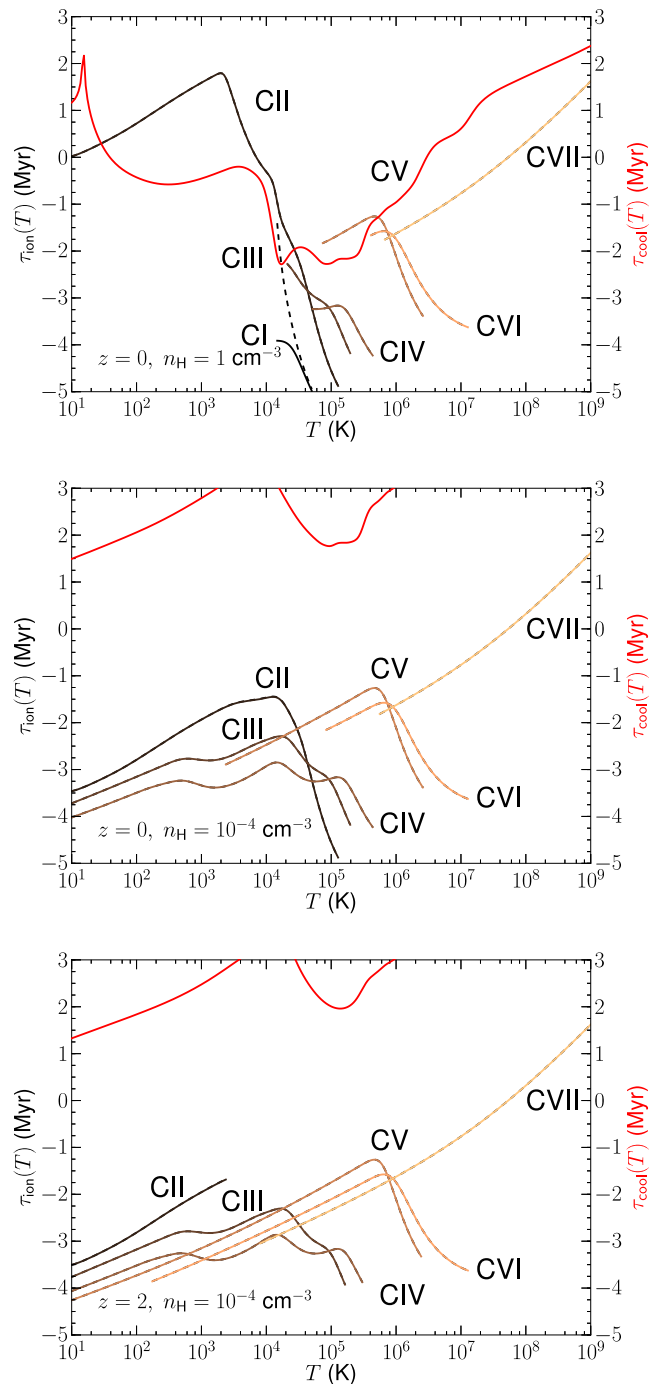


Figure 2. Carbon ions ionization time-scale, τ_{ion} (coloured curves), and cooling time-scale, τ_{cool} (red curve), as a function of temperature for different redshifts and densities (as indicated in the panels). τ_{ion} is only plotted if the fraction of the corresponding C ion is above 0.001. The dashed curves trace the ionization time-scale for pure collisional ionization (this only makes a difference for C I at high densities).

3 IONIZATION BALANCE

3.1 Density dependence

In Fig. 3, we show the ionization equilibrium (i.e. the fraction of the atoms of a given element that come in the form of a given ion) of hydrogen (H), oxygen (O), carbon (C) and iron (Fe) as a function of temperature, calculated for the UVB at redshift $z = 0$ and a density of

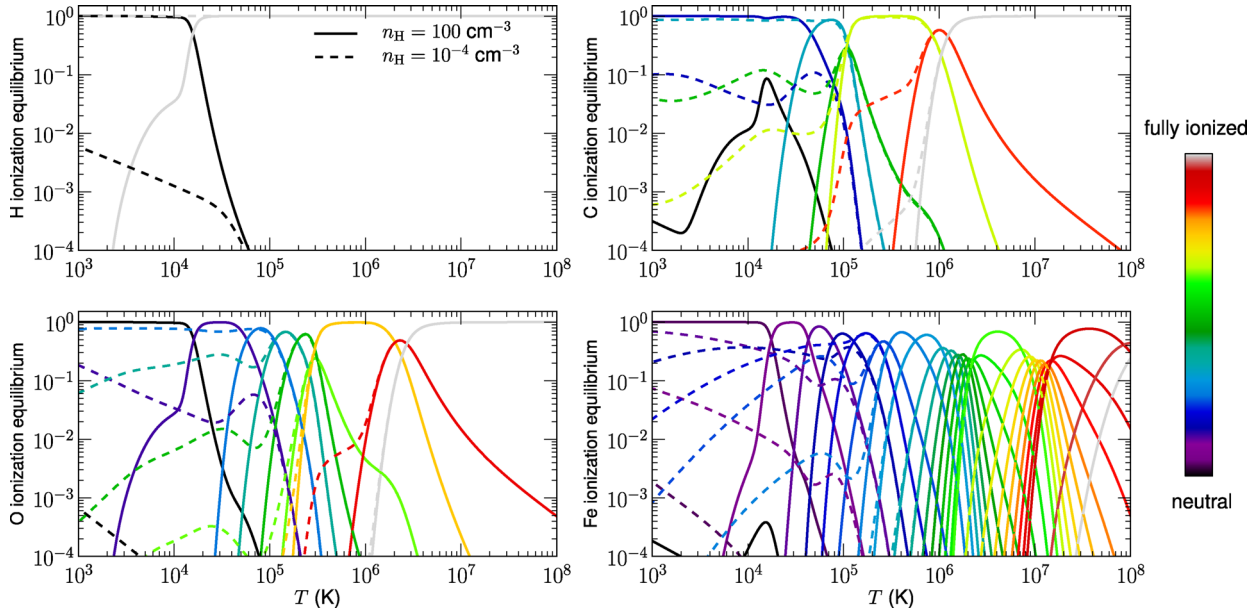


Figure 3. Ionization equilibrium of hydrogen (H), oxygen (O), carbon (C) and iron (Fe) as a function of temperature, calculated for the UVB at redshift $z = 0$ and a density of $n_{\text{H}} = 100 \text{ cm}^{-3}$ (full lines) and $n_{\text{H}} = 10^{-4} \text{ cm}^{-3}$ (dashed lines). The colour scale indicates the ionization stage of the various ions.

$n_{\text{H}} = 100 \text{ cm}^{-3}$ (full lines) and $n_{\text{H}} = 10^{-4} \text{ cm}^{-3}$ (dashed lines). At densities above the self-shielding density threshold, hydrogen can recombine at temperatures below about 20 000 K, thus shielding the gas from the photo-ionizing UVB. Likewise, oxygen, with its first ionization potential very close to that of hydrogen, recombines to its neutral form. Only elements with first ionization potentials smaller than that of hydrogen, such as carbon and iron in this example, remain once ionized below $\sim 10\,000$ K.

At low densities, below the self-shielding density threshold of hydrogen, the ionizing UVB can flood the gas unimpeded, keeping over 99 per cent of the hydrogen ionized. This completely erases the contribution of hydrogen to the cooling via its free-bound and bound-bound transitions. In this example, the UVB affects

essentially all ions of carbon and oxygen, with C III , O III and Fe IV the most common ionization stages of these elements at low temperatures. The high abundance of free photo-electrons provides extra cooling through radiative free-bound transitions, more than making up for the lack of efficient coolants such as the C II and Fe II infrared fine-structure lines at low temperatures (see below, in Section 4).

3.2 Redshift dependence

For densities below the hydrogen self-shielding density threshold, the ionizing strength of the UVB, which varies significantly with redshift, has a profound impact on the ionization balance. In Fig. 4,

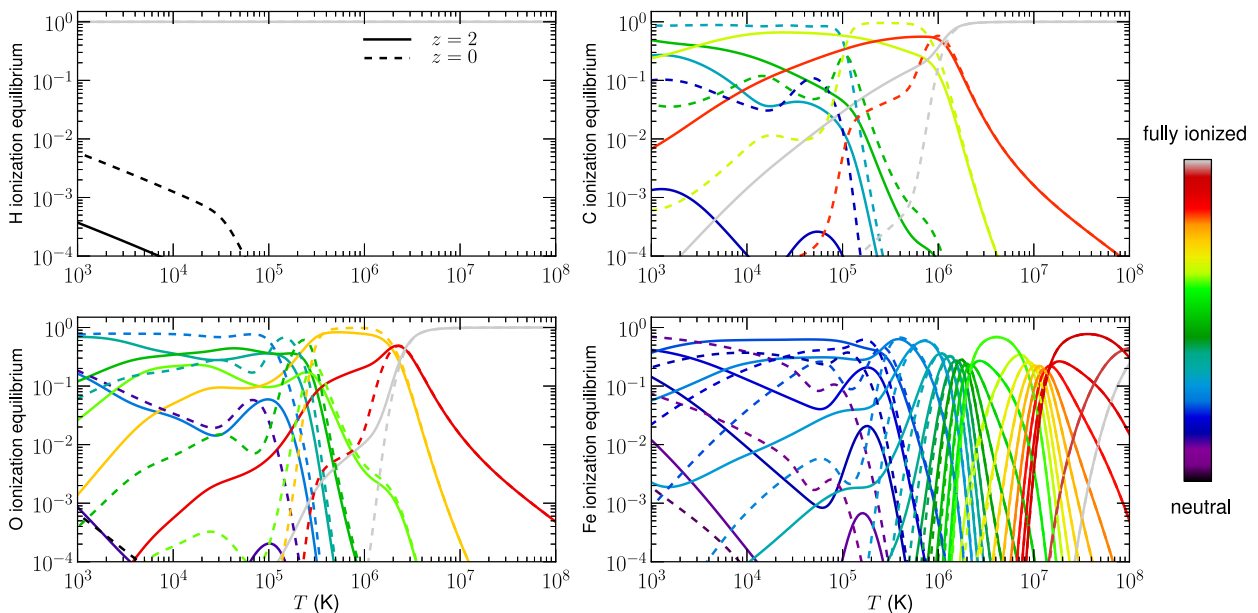


Figure 4. Ionization equilibrium of hydrogen (H), oxygen (O), carbon (C) and iron (Fe) as a function of temperature, calculated for a density of $n_{\text{H}} = 10^{-4} \text{ cm}^{-3}$ for a $z = 2$ UVB (full lines) and $z = 0$ UVB (dashed lines). The colour scale indicates the ionization stage of the various ions.

we show the ionization equilibrium of hydrogen (H), oxygen (O), carbon (C) and iron (Fe) as a function of temperature, calculated for a gas with a density of $n_{\text{H}} = 10^{-4} \text{ cm}^{-3}$ and subjected to a $z = 2$ UVB (full lines) and a $z = 0$ UVB (dashed lines). Clearly, the stronger $z = 2$ UVB leads to more ionization of the various elements with C III and C IV the dominant carbon ions at low temperatures while oxygen is found predominately in its O IV and O V forms. Below 10^5 K iron exists mostly as Fe VII and Fe VIII .

4 COOLING

4.1 Density dependence

In the top panel of Fig. 5, we compare cooling curves, denoted by $\Lambda(T)$, calculated using the $z = 0$ UVB, $[\text{Fe}/\text{H}] = 0$, and for different densities between $n_{\text{H}} = 10^{-9} \text{ cm}^{-3}$ and $n_{\text{H}} = 100 \text{ cm}^{-3}$, as indicated in the figure.

At the lowest densities, inverse Compton scattering dominates the cooling rate at all temperatures and the cooling curve is virtually featureless. At higher densities, in the range $10^{-9} < n_{\text{H}} \ll 10^{-2} \text{ cm}^{-3}$, cooling via radiative free-bound and bound-bound transitions becomes important at temperatures below $\sim 10^7 \text{ K}$. At these densities, hydrogen is essentially completely photo-ionized by the UVB and does not contribute to the cooling rate via these transitions. The many free electrons keep the rate of free-bound transitions high even at the lowest temperatures considered here. At the highest temperatures, upwards of $\sim 10^7 \text{ K}$, the cooling curve's temperature slowly changes from the $\Lambda \propto T$ behaviour of inverse Compton scattering to the less steep $\Lambda \propto \sqrt{T}$ behaviour that is characteristic of free-free transitions. At the highest densities, for $n_{\text{H}} \gtrsim 10^{-2} \text{ cm}^{-3}$, hydrogen is able to recombine and shield the other elements from the UVB. Consequently, the 10 000–20 000 K peak in the cooling rate contributed by free-bound and bound-bound transitions of hydrogen appears. As a result of the strong reduction of the free electron density, the cooling rate is likewise strongly reduced below $T \sim 100 \text{ K}$. Below 10^4 K , cooling via infrared finestructure emission lines, predominantly the $157.7 \mu\text{m}$ line of C II , dominates.

4.2 Redshift dependence

The cosmological redshift enters the picture in two distinct ways: on the one hand, it controls the strength of the UVB, and on the other, it determines the contribution to the cooling rate from inverse Compton scattering through its $(1+z)^4$ -dependence. For the zero-strength $z = 15$ UVB, H always fully recombines below $T \sim 10^4 \text{ K}$ and the cooling rate plummets by 4 orders of magnitude, as is evident in the bottom panel of Fig. 5. The only contribution to the inverse Compton scattering cooling rate now comes from the handful of ionization electrons from ions with small ionizing potentials, such as C I , Fe I , etc.

The stronger the UVB, the more highly ionized the different elements are. This lack of lowly ionized species leads to a dramatic decrease in the cooling rate due to free-bound and bound-bound emission (compare e.g. the cooling curves for densities $n_{\text{H}} = 10^{-6} \rightarrow 10^{-4} \text{ cm}^{-3}$ between $z = 0$ and $z = 2$).

4.3 $[\text{Fe}/\text{H}]$ dependence

In Fig. 6, we compare the cooling rates at metallicities between $[\text{Fe}/\text{H}] = -\infty$ and $[\text{Fe}/\text{H}] = 0.5$ (as indicated in the figure), cal-

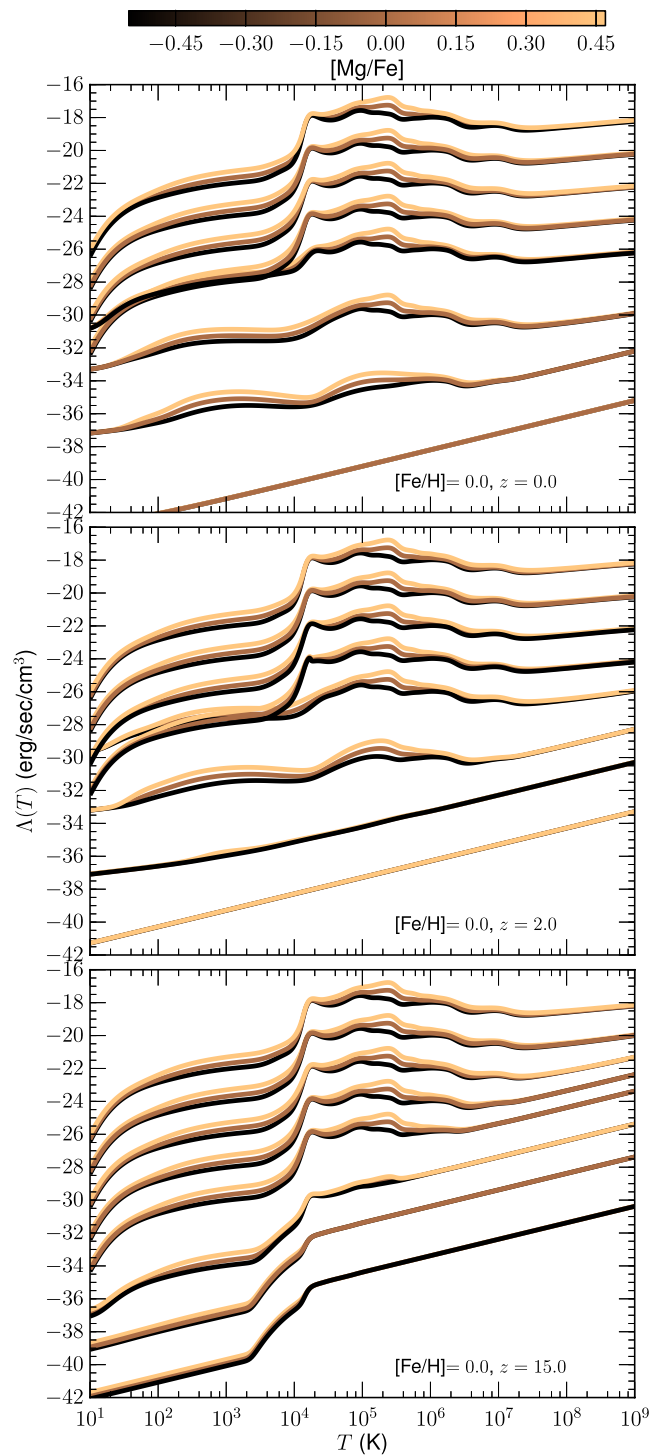


Figure 5. Logarithm of the cooling rate as a function of temperature for $[\text{Fe}/\text{H}] = 0.0$, and for the $z = 0$ (top panel), $z = 2$ (middle panel) and $z = 15$ (bottom panel) cosmic UVB, plotted for different densities (from the top curve downwards: $n_{\text{H}} = 100, 10, 1, 10^{-1}, 10^{-2}, 10^{-4}, 10^{-6}, 10^{-9} \text{ cm}^{-3}$) and $[\text{Mg}/\text{Fe}]$ -values (colour code).

culated for Mg-abundance $[\text{Mg}/\text{Fe}] = 0.0$, density $n_{\text{H}} = 0.01 \text{ cm}^{-3}$ and for redshifts $z = 0, z = 8$ and $z = 15$.

At high temperatures, inverse Compton scattering and free-free interactions dominate the cooling rate at high redshift. In this regime, the metallicity only affects the density of free electrons

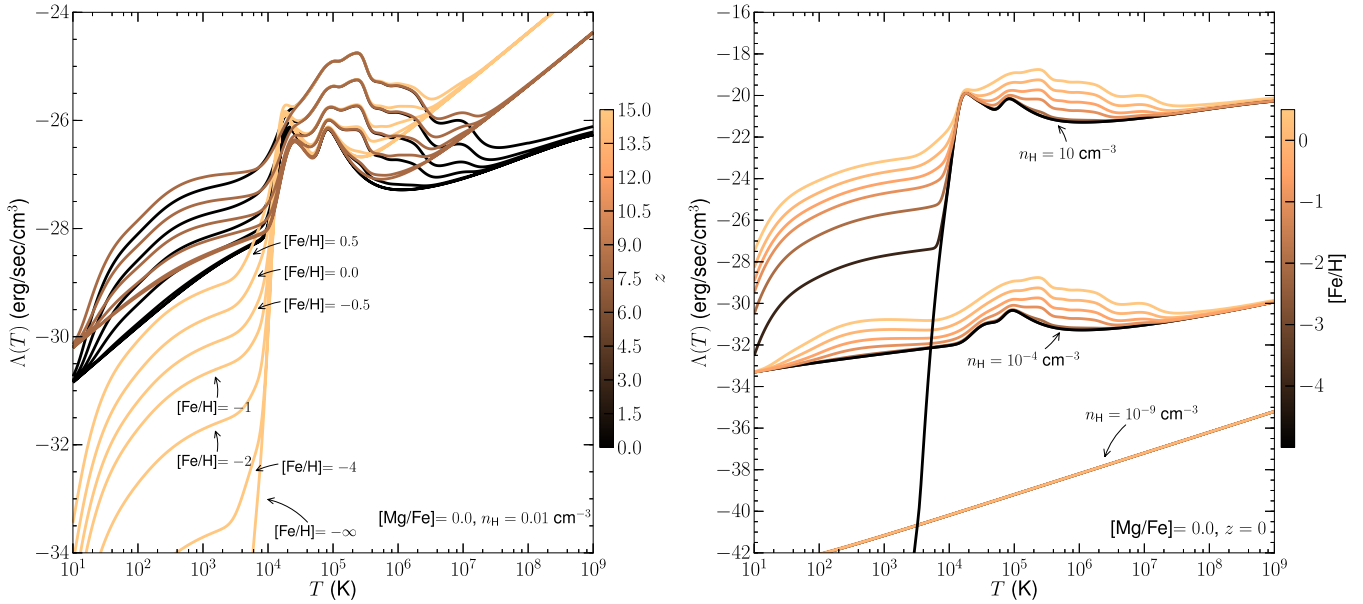


Figure 6. Left-hand panel: logarithm of the cooling rate as a function of temperature for Mg-abundance $[Mg/Fe]=0.0$, density $n_H = 0.01 \text{ cm}^{-3}$, for redshifts $z = 0$, $z = 8$ and $z = 15$ (colour code), and for metallicities between $[Fe/H] = -\infty$ and $[Fe/H] = 0.5$ (as indicated in the figure). Right-hand panel: the logarithm of the cooling rate as a function of temperature for Mg-abundance $[Mg/Fe] = 0.0$, redshift $z = 0$, densities $n_H = 10^{-9} \text{ cm}^{-3}$, $n_H = 10^{-4} \text{ cm}^{-3}$ and $n_H = 10 \text{ cm}^{-3}$, and for different metallicities (colour code). The bottom, black curve of each series is calculated for $[Fe/H] = -\infty$.

and, as a result, the cooling curves are not very sensitive to metallicity. At temperatures $T \lesssim 10^7 \text{ K}$, partially ionized atoms can exist and free-bound and bound-bound transitions contribute greatly to the cooling rate. Likewise, for $T \lesssim 10^4 \text{ K}$, the cooling rate via infrared fine-structure lines is a strong function of metallicity since in this regime the cooling depends crucially on the presence of a few key ions. This is especially true when the UVB is very weak or even absent (as at $z = 15$). In that case, hydrogen fully recombines below $T \sim 10^4 \text{ K}$, consuming all free electrons, and the cooling rate drops sharply. The cooling contributed by infrared fine-structure lines of metal ions can then make a huge difference.

4.4 [Mg/Fe] dependence

In Fig. 5, the cooling curves are colour coded according to their $[Mg/Fe]$ -value. The amount of α -enhancement clearly has a great impact on the cooling rate in those temperature ranges where key ions of the α -elements contribute free-bound and bound-bound cooling. Around $T \sim 200\,000 \text{ K}$, depending on the abundance of O and Ne, the cooling rate can vary by up to an order or magnitude. Around $T \sim 10^6 \text{ K}$, Si lets its presence be felt. In the range $10^2 \lesssim T \lesssim 10^4 \text{ K}$, infrared fine-structure emission lines from α -element ions such as $O\text{I}$ and $Si\text{II}$ contribute to the cooling rate (Maio et al. 2007) and can make an order of magnitude difference depending on whether a gas parcel has been enriched only by the SNIa and low-mass stars (the ‘slow’ contribution to the yield) or only by SNIi and massive stars (the ‘fast’ contribution to the yield). Since the ‘fast’ and ‘slow’ yields of carbon happen to be very similar, the cooling rate is relatively insensitive to $[Mg/Fe]$ below 100 K . For a very weak or absent UVB, ions of α -elements with small first ionization potentials, ionized by the ISRF, contribute free electrons and slightly raise the cooling rate below $T \sim 10^4 \text{ K}$ (this is most noticeable in the bottom panel of Fig. 5).

5 HEATING

The heating rate is a very strong function of gas density and metallicity. Both parameters determine the density of partially ionized atoms capable of absorbing energy from the UVB through further ionization while the former, moreover, sets the amount of self-shielding. At low densities, the heating rate is essentially a monotonically declining function of temperature: the higher the temperature the lower the densities of the lowly ionized species that absorb heat most efficiently. The heating rate also increases with increasing metallicity since this obviously raises the number of heat absorbing ions.

At high densities, the heating rate shows a much more complex behaviour. While the densities of near-neutral species increase towards lower temperatures, potentially raising the heat-absorbing capabilities of the gas, the self-shielding by neutral hydrogen suppresses the UVB. This can lead to a plateau in the heating rate for temperatures below 10^4 K , see e.g. at $n_H = 0.1 \text{ cm}^{-3}$ in Fig. 7. This plateau is simply the heating rate of a fully neutral hydrogen gas irradiated by a strongly reduced UVB. For higher metallicities, lowly ionized metals provide a small amount of extra heating above this plateau.

At the highest densities, the hydrogen-ionizing part of the UVB is almost completely suppressed and other sources of heating, such as the ISRF, become significant. Since hydrogen cannot extract heat from the ISRF employed in our calculations, the heating rate drops to zero for $[Fe/H] = -\infty$ when hydrogen recombines. For non-zero metallicities, the ISRF can ionize and thus heat those elements that have ionization potentials smaller than that of hydrogen, explaining the complex behaviour of the heating rate between 10^3 and 10^4 K .

The $[Mg/Fe]$ -dependence of the heating rates is much weaker than their metallicity dependence. As an example, we show the heating rate as a function of temperature for Fe-abundance $[Fe/H] = 0.0$, redshift $z = 0$, gas density $n_H = 100 \text{ cm}^{-3}$ and different $[Mg/Fe]$ ratios in Fig. 8. While $[Mg/Fe]$ varies with one dex, the heating rate changes by a factor of 5 at most.

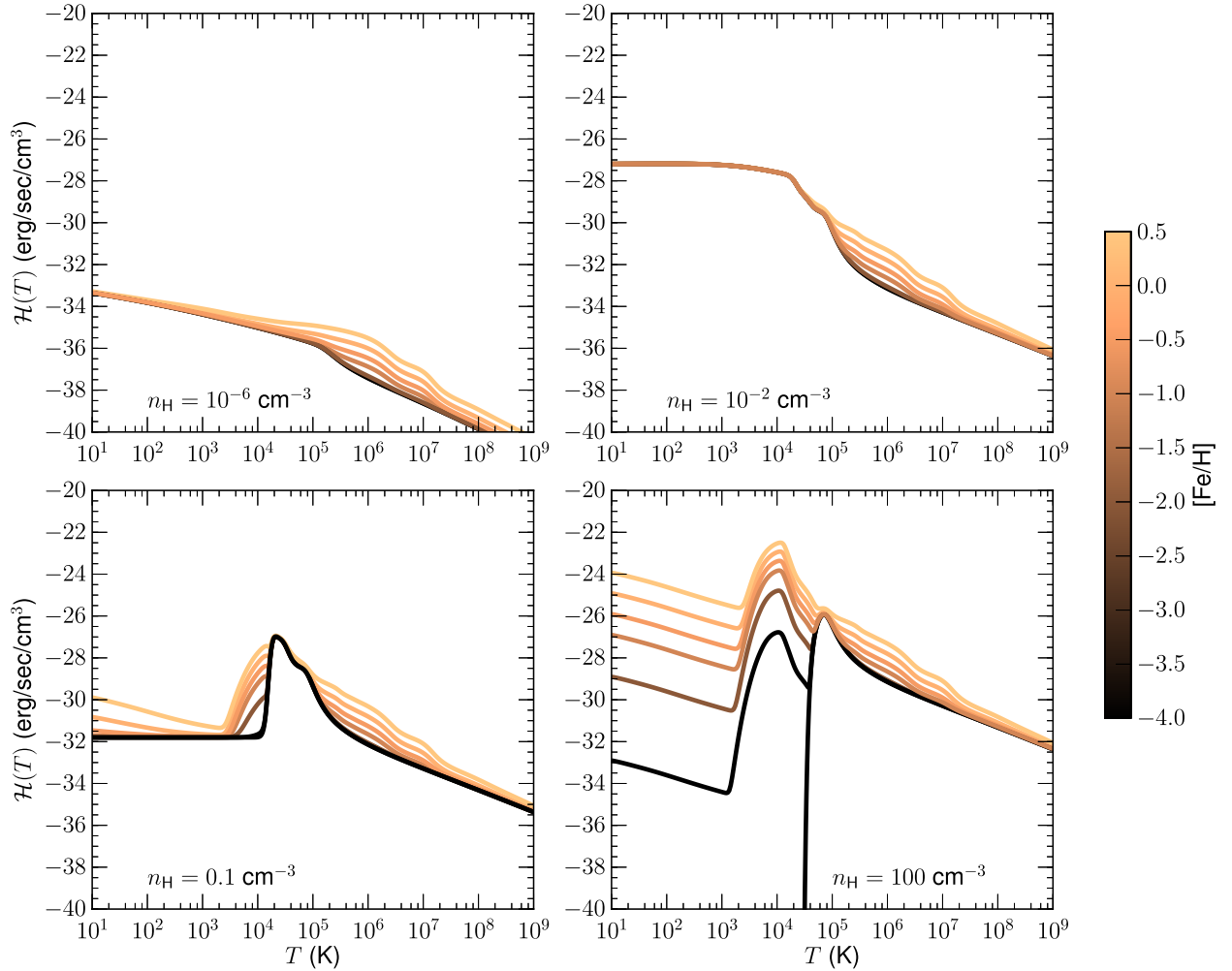


Figure 7. Logarithm of the heating rate as a function of temperature for Mg-abundance $[\text{Mg}/\text{Fe}] = 0.0$, redshift $z = 0$, Fe-abundance $[\text{Fe}/\text{H}]$ (colour code) and different gas densities (indicated in the panels).

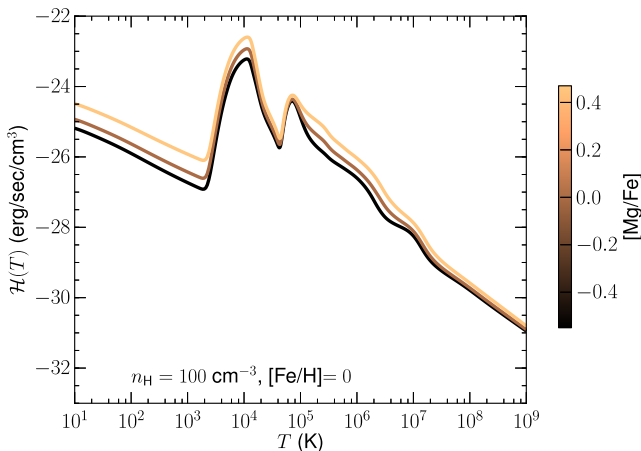


Figure 8. Logarithm of the heating rate as a function of temperature for Fe-abundance $[\text{Fe}/\text{H}] = 0.0$, redshift $z = 0$, gas density $n_{\text{H}} = 100 \text{ cm}^{-3}$ and different $[\text{Mg}/\text{Fe}]$ ratios (colour code).

In Fig. 9, we plot the net cooling rates, defined as $\Lambda_{\text{net}} = |\Lambda - \mathcal{H}|$, of solar metallicity gas exposed to the $z = 0$ (top panel), $z = 2$ (middle panel) and $z = 8$ (bottom panel) UVB for different densities, with self-shielding (full lines) and without self-shielding

(dotted lines). Clearly, for densities below the self-shielding threshold density of $n_{\text{H1,crit}} = 0.007 \text{ cm}^{-3}$ switching shielding on or off makes no difference: the UVB can fully penetrate, ionize and heat the gas. Above the self-shielding threshold (roughly the top four curves in each panel), the gas is dense enough to recombine and, with self-shielding switched on, to strongly suppress the UVB. As a consequence, the heating rate drops steeply below $T \sim 10^4 \text{ K}$. With self-shielding switched off, hydrogen also largely recombines but now does not suppress the UVB. This has a profound influence on the heating rate which keeps on increasing towards lower temperatures until reaching a plateau below $T \sim 10^4 \text{ K}$, as discussed earlier. Without self-shielding against the UV radiation, ionization levels tend to be higher, leading to higher abundances of important cooling ions such as C II and, consequently, to higher cooling rates in the $T < 10^4 \text{ K}$ temperature regime. Self-shielding also has a strong effect on the equilibrium temperature of the gas. If, without self-shielding, the UVB can flood the gas unimpeded, the equilibrium temperature can be over two orders of magnitude higher than in the self-shielded case (e.g. $T_{\text{eq}} = 6500 \text{ K}$ without self-shielding versus $T_{\text{eq}} = 20 \text{ K}$ with self-shielding for the $z = 2$ UVB at gas density $n_{\text{H1}} = 0.1 \text{ cm}^{-3}$).

The most striking effect of the self-shielding prescription is the equilibrium temperature to which the gas would evolve, given sufficient time. The sudden downward break in each curve Λ_{net}

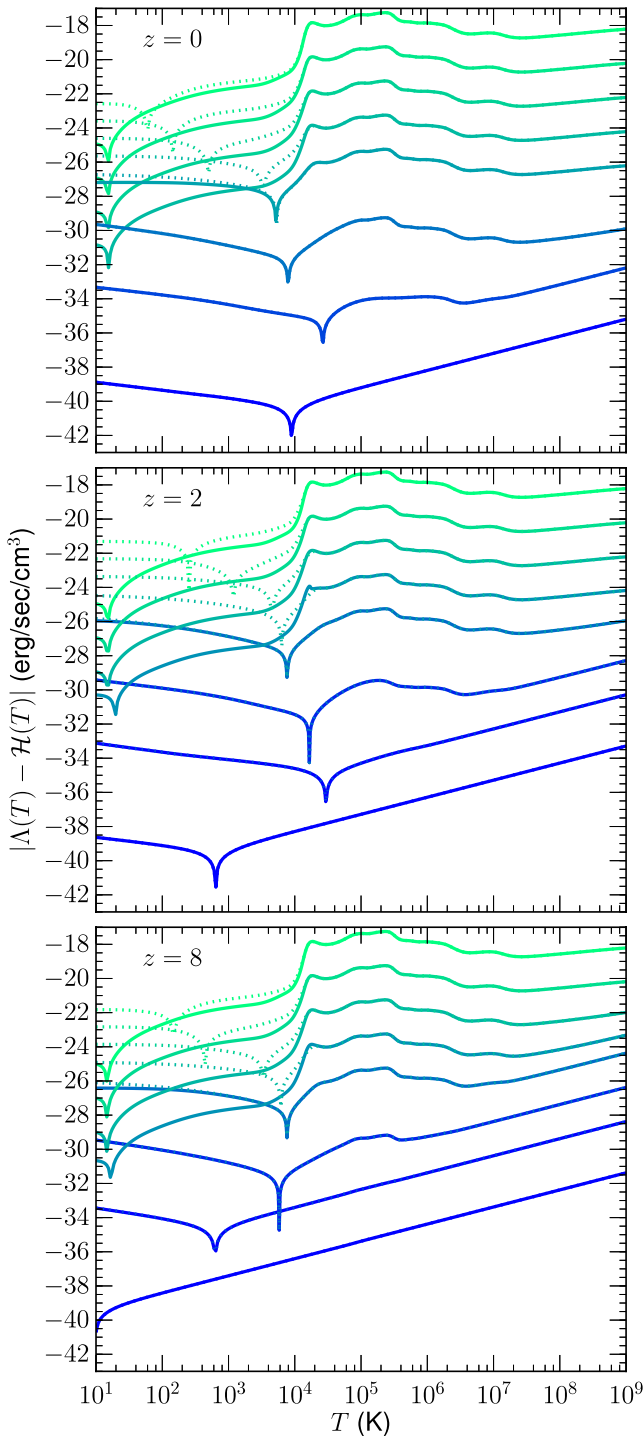


Figure 9. Logarithm of the net cooling rate of solar metallicity gas exposed to the $z = 0$ UVB (top panel), the $z = 2$ UVB (middle panel) and the weak $z = 8$ UVB (bottom panel) with self-shielding (full lines) and without self-shielding (dotted lines). The line colours reflect gas density (from the top curve downwards: $n_{\text{H}} = 100, 10, 1, 10^{-1}, 10^{-2}, 10^{-4}, 10^{-6}$ and 10^{-9} cm^{-3}).

indicates where $\Lambda = \mathcal{H}$ and hence marks the equilibrium temperature. In the middle panel of Fig. 9, the strong $z = 2$ UVB is employed while in the top and bottom panels, the results for the much weaker $z = 0$ and $z = 8$ UVBs are shown. Without self-shielding, gas is incapable of radiatively cooling significantly below $T \sim 10^4 \text{ K}$, except

for sufficiently high densities and at large enough redshifts where the UVB is still weak. Except at late and at very early cosmic times, the UVB severely inhibits the formation of the cold, high-density clouds in which stars are thought to form while it facilitates the removal of low-density gas by ram-pressure stripping and galaxy interactions (Mayer et al. 2007; Governato et al. 2010). Together with supernova feedback, which most strongly affects high-density regions, this almost completely extinguishes star formation in simulated dwarf galaxies after $z \sim 2$ (Simpson et al. 2012). However, while the specific star formation rate of Local Volume dwarf galaxies was generally larger before $z \sim 2$ compared with later times, they do show a wide variety of more or less continuous star formation histories over their full lifetimes (Weisz et al. 2011). It remains to be seen whether a proper, self-consistent treatment of the effects of photo-heating by the UVB, including the effects of self-shielding, on the ionization equilibrium and the resultant cooling and heating rates, can alleviate this problem.

Moreover, the non-inclusion of self-shielding makes it virtually impossible to clearly identify different phases in the ISM of a simulated galaxy and to compare them with observed galaxies. For instance, in their analysis of a fully cosmological hydrodynamical dwarf galaxy which was simulated including a UVB but neglecting self-shielding, Pilkington et al. (2011) find that gas particles typically have temperatures of the order of 7000–9000 K. In order to investigate the simulated dwarf’s H I properties, these authors are forced to select as ‘cold’ gas particles those with temperatures below 15 000 K. Since the calculation of the cooling and heating curves presented here involves the determination of the ionization equilibrium as a function of the gas properties, we immediately have the neutral hydrogen fraction at our disposal. As can be seen in Fig. 10, in the presence of a UVB, the neutral fraction of a gas parcel is not only temperature dependent but also strongly density and redshift dependent. The black dots in this figure indicate the neutral fraction at the equilibrium temperature for each redshift and gas density. Clearly, the hydrogen gas has to be denser than several 0.01 cm^{-3} before it can become neutral (i.e. $n_{\text{HI}}/n_{\text{H}} \gtrsim 0.95$). This broadly agrees with the estimated lower bound on the density of the Warm Neutral Medium in the Milky Way (Wolfire et al. 2003).

6 DISCUSSION

In this paper, we have presented a new calculation of composition-dependent cooling and heating curves intended for use in numerical simulations of galaxy formation and evolution. For each elemental mix, density, temperature and cosmological time the ionization equilibrium was determined using a modified version of CHIANTI, equipped with collisional and radiative ionization, by the cosmic UVB and an ISRF, and charge-transfer reactions. We believe these curves address several drawbacks of currently available tabulations of cooling rates.

We have shown that the full range of abundance variations likely to be encountered in stars and neutral and ionized gas in a galaxy can be described adequately by a simplified chemical-evolution model in which there are only two contributions to the elemental yields: a ‘fast’ one (encompassing the contributions from SNII and massive IMS) and a ‘slow’ one (encompassing the contributions from SNIa and less massive IMS). The ratio of both contributions can be linked directly to the Fe and Mg abundances which provide us with two strong handles on *all* other element abundances.

Thus, the cooling and heating curves depend on only five parameters (temperature, density, redshift, [Fe/H] and [Mg/Fe]). They are easily tabulated, and can be efficiently interpolated during a

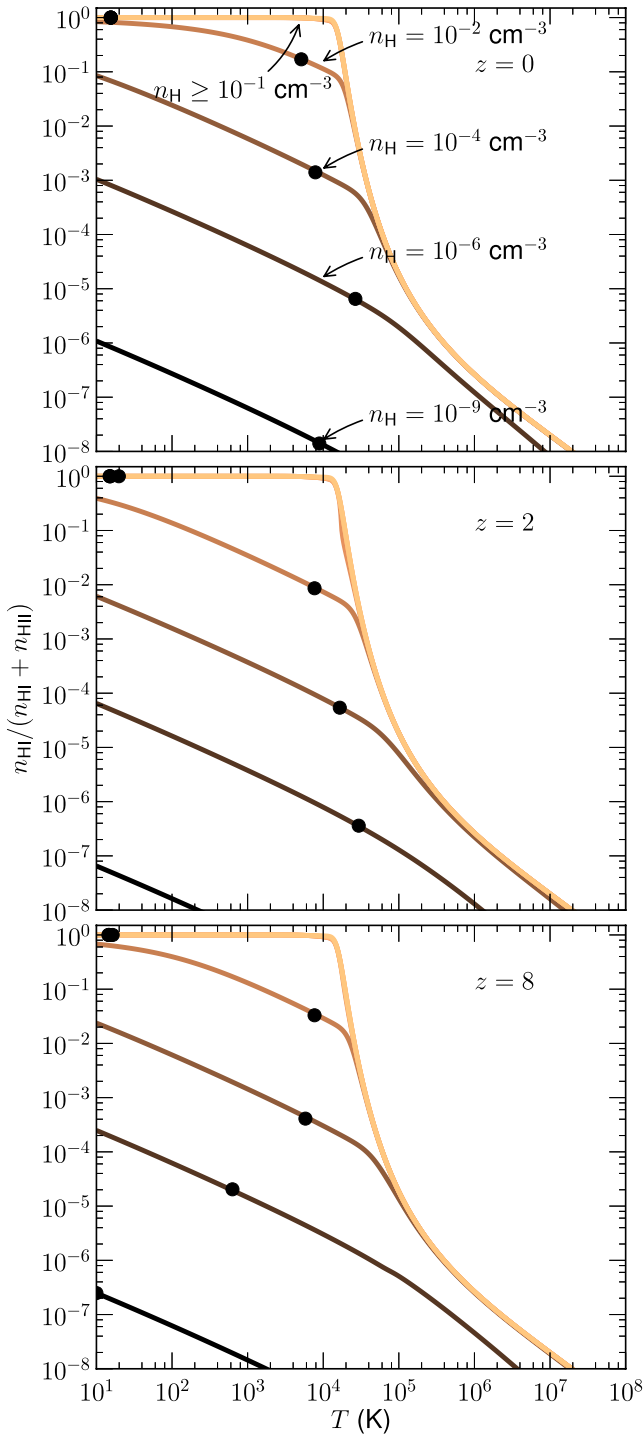


Figure 10. Neutral fraction of solar metallicity gas exposed to the $z = 0$ UVB (top panel), the $z = 2$ UVB (middle panel) and the weak $z = 8$ UVB (bottom panel), as a function of temperature and density (indicated in the figure, from the top curve downwards: $n_{\text{H}} = 100, 10, 1, 10^{-1}, 10^{-2}, 10^{-4}, 10^{-6}$ and 10^{-9} cm^{-3}). The black dots indicate the neutral fraction at the equilibrium temperature for each density and redshift.

simulation. We have implemented a five-dimensional interpolator in our own simulation code enabling us to employ the cooling and heating curves presented here in galaxy evolution simulations. A detailed analysis of the effects of using these new curves on such

simulations, especially regarding the evolution of the star formation rate in dwarf galaxies, will be presented in a forthcoming paper.

We believe that the main advantage of this work is that, within the confines of a well-defined chemical evolution model and adopting the ionization equilibrium approximation, it provides accurate cooling and heating curves for a wide range of physical and chemical gas properties. These should be valid as long as the gas is neutral or, partially, ionized (molecule and dust formation have not been taken into account). Moreover, during a numerical simulation, one need only follow the evolution of the Fe and Mg abundances, leading to a reduction in memory requirements and computational cost compared to the element-by-element approach. It takes into account self-shielding in an approximate but motivated and well-tested way. This is key to resolving the formation of cold, neutral, high-density clouds suitable for star formation and to studying the structure of the multi-phase ISM in galaxy simulations. Moreover, since we have stored the ionization equilibrium for each combination of temperature, density, redshift, $[\text{Fe}/\text{H}]$ and $[\text{Mg}/\text{Fe}]$, we can in principle calculate any desired physical property of the gas using `CHIANTI.PY`.

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Pre-compiled tables of the cooling and heating curves are available to the community as online-only supporting information. These tables, future updates and new material can also be found on and downloaded from <http://users.ugent.be/~sdrijcke>.

REFERENCES

- Aubert D., Teyssier T., 2010, *ApJ*, 724, 244
 Bekki K., 2013, preprint (arXiv:1304.1633)
 Brook C. B., Stinson G., Gibson B. K., Roškar R., Wadsley J., Quinn T., 2012, *MNRAS*, 419, 771
 Chabrier G., 2003, *ApJ*, 586, L133
 Cloet-Osselaer A., De Rijcke S., Schroyen J., Dury V., 2012, *MNRAS*, 423, 735
 Dere K. P., Landi E., Young P. R., Del Zanna G., Landini M., Mason H. E., 2009, *A&A*, 498, 915
 Faucher-Giguère C.-A., Lidz A., Zaldarriaga M., Hernquist L., 2009, *ApJ*, 703, 1416
 Faucher-Giguère C.-A., Kereš D., Dijkstra M., Hernquist L., Zaldarriaga M., 2010, *ApJ*, 725, 633
 Fitzpatrick E. L., 2010, *ApJ*, 725, 2401
 François P., Matteucci F., Cayrel R., Spite M., Spite F., Chiappini C., 2004, *A&A*, 421, 613
 Gabor J. M., Davé R., 2012, *MNRAS*, 427, 1816
 Gavilán M., Buell J. F., Mollá M., 2005, *A&A*, 432, 861
 Gnat O., Sternberg A., 2007, *ApJS*, 168, 213
 Gnedin N. Y., Hollon N., 2012, *ApJS*, 202, 13
 Governato F. et al., 2010, *Nat*, 463, 203
 Grevesse N., Asplund M., Sauval A. J., Scott P., 2010, *Astrophys. Space Sci.*, 328, 179
 Kim J.-H., Krumholz M. R., Wise J. H., Turk M. J., Goldbaum N. J., Abel T., 2012, preprint (arXiv:1210.3361)
 Kingdon J. B., Ferland G. J., 1996, *ApJS*, 106, 205
 Krumholz M. R., 2012, *ApJ*, 759, 9
 Landstreed J. D., 2011, *A&A*, 528, A132
 Maio U., Dolag K., Ciardi B., Tornatore L., 2007, *MNRAS*, 379, 963
 Mathis J. S., Mezger P. G., Panagia N., 1983, *A&A*, 128, 212
 Mayer L., Kazantzidis S., Mastropietro C., Wadsley J., 2007, *Nat*, 445, 738
 Nagamine K., Choi J., Yajima H., 2010, *ApJ*, 725, L219

- Nomoto K., Hashimoto M., Tsujimoto T., Thielemann F.-K., Kishimoto N., Kubo Y., Nakasato N., 1997, *Nucl. Phys.*, 616, 79
- Pilkington K. et al., 2011, *MNRAS*, 417, 2891
- Revaz Y. et al., 2009, *A&A*, 501, 189
- Sawala T., Guo Q., Scannapieco C., Jenkins A., White S., 2011, *MNRAS*, 413, 659
- Scannapieco C., White S. D. M., Springel V., Tissera P. B., 2011, *MNRAS*, 417, 154
- Schaye J. et al., 2010, *MNRAS*, 402, 1536
- Schroyen J., De Rijcke S., Valcke S., Cloet-Osselaer A., Dejonghe H., 2011, *MNRAS*, 416, 601
- Shen S., Wadsley J., Stinson G., 2010, *MNRAS*, 407, 1581
- Simpson C. M., Bryan G. L., Johnston K. V., Smith B. D., Mac Low M.-M., Sharma S., Tumlinson J., 2012, preprint (arXiv:1211.1071v1)
- Smith B., Sigurdsson S., Abel T., 2008, *MNRAS*, 385, 1443
- Smith B. D., Hallman E. J., Shull J. M., O'Shea B. W., 2011, *ApJ*, 731, 6
- Stancil P. C., Schultz D. R., Kimura M., Gu J.-P., Hirsch G., Buenker R. J., 1999, *A&AS*, 140, 225
- Stancil P. C. et al., 1998, *ApJ*, 502, 1006
- Sutherland R. S., Dopita M. A., 1993, *ApJS*, 88, 253
- Tajiri Y., Umemura M., 1998, *ApJ*, 502, 59
- Tepper-García T., Richter P., Schaye J., Booth C. M., Dalla Vecchia C., Theuns T., Wiersma R. P. C., 2011, *MNRAS*, 413, 190
- Tolstoy E., Hill V., Tosi M., 2009, *ARA&A*, 47, 371
- Tsujimoto T., Nomoto K., Yoshii Y., Hashimoto M., Yanagida S., Thielemann F.-K., 1995, *MNRAS*, 277, 945
- Vasiliev E. O., 2013, *MNRAS*, 431, 638
- Verner D. A., Ferland G. J., Korista K. T., Yakovlev D. G., 1996, *ApJ*, 465, 487
- Weisz D. R. et al., 2011, *ApJ*, 739, 27
- Wiersma R. P. C., Schaye J., Smith B. D., 2009, *MNRAS*, 393, 99
- Wolfire M. G., McKee C. F., Hollenbach D., Tielens A. G. G. M., 2003, *ApJ*, 587, 278
- Worley C. C., Cottrell P. L., Freeman K. C., Wylie-de Boer E. C., 2009, *MNRAS*, 400, 1039
- Worthey g., Faber S. M., Gonzalez J. J., 1992, *ApJ*, 398, 69
- Yajima H., Jun-Hwan C., Nagamine K., 2011, *MNRAS*, 412, 411

SUPPORTING INFORMATION

Additional Supporting Information may be found in the online version of this article:

Pre-compiled tables of the cooling and heating curves (<http://mnras.oxfordjournals.org/lookup/suppl/doi:10.1093/mnras/stt942/-/DC1>).

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