






ELUCID. VII. Using Constrained Hydro Simulations to Explore the Gas Component of the Cosmic Web

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Abstract

Using reconstructed initial conditions in the Sloan Digital Sky Survey (SDSS) survey volume, we carry out constrained hydrodynamic simulations in three regions representing different types of the cosmic web: the Coma cluster of galaxies; the SDSS Great Wall; and a large low-density region at $z \sim 0.05$. These simulations, which include star formation and stellar feedback but no active galactic nucleus formation and feedback, are used to investigate the properties and evolution of intergalactic and intracluster media. About half of the warm-hot intergalactic gas is associated with filaments in the local cosmic web. Gas in the outskirts of massive filaments and halos can be heated significantly by accretion shocks generated by mergers of filaments and halos, respectively, and there is a tight correlation between the gas temperature and the strength of the local tidal field. The simulations also predict some discontinuities associated with shock fronts and contact edges, which can be tested using observations of the thermal Sunyaev–Zel'dovich effect and X-rays. A large fraction of the sky is covered by Ly α and O VI absorption systems, and most of the O VI systems and low-column-density H I systems are associated with filaments in the cosmic web. The constrained simulations, which follow the formation and heating history of the observed cosmic web, provide an important avenue to interpret observational data. With full information about the origin and location of the cosmic gas to be observed, such simulations can also be used to develop observational strategies.

Unified Astronomy Thesaurus concepts: [Hydrodynamical simulations \(767\)](#); [Intergalactic medium \(813\)](#); [Intracluster medium \(858\)](#); [Intergalactic filaments \(811\)](#); [Sunyaev-Zeldovich effect \(1654\)](#); [X-ray photometry \(1820\)](#); [Quasar absorption line spectroscopy \(1317\)](#); [Coma Cluster \(270\)](#); [Cosmic web \(330\)](#); [Voids \(1779\)](#); [Large-scale structure of the universe \(902\)](#); [Warm-hot intergalactic medium \(1786\)](#)

1. Introduction

In the local universe, only about 10% of the baryons are found to be locked in stars and dense interstellar media (e.g., Bregman 2007). Most of the baryons are, therefore, expected to be ionized and diffuse in various components of the cosmic web, from virialized DM halos to large-scale filaments and sheets. The thermal, dynamical, and chemical states of the baryonic gas are subject to a number of processes, such as gravitational collapse, radiative heating–cooling, and feedback from stars and active galactic nuclei (AGNs), all closely related to the formation and evolution of the objects we observe. Clearly, the study of the gas components in the cosmic web is a key step to understanding the formation and evolution of galaxies and the large-scale structure of the universe.

Observationally, the gas components can be investigated using various methods, such as X-ray emission, Sunyaev–Zel'dovich (SZ) effects, and emission and absorption lines. The intracluster

medium (ICM), which is the hottest component in the cosmic web, can be observed as extended X-ray sources (e.g., Mirakhor & Walker 2020; Churazov et al. 2021). The free electrons in the hot ICM can also change the energy distribution of the cosmic microwave background photons via inverse Compton scattering, producing a thermal SZ effect (tSZ; Sunyaev & Zel'dovich 1972) that can be used to investigate the thermal content of the ICM (e.g., Planck Collaboration et al. 2013a; Bleem et al. 2015). Since the X-ray emission and tSZ effect depend on gas density and temperature in different ways, the combination of X-ray and tSZ observations can provide measurements that can be used to infer the thermodynamic properties of the ICM (e.g., Ghirardini et al. 2018; Mirakhor & Walker 2020; Churazov et al. 2021). The gas in smaller and more abundant groups is colder and thus more difficult to detect individually. A common practice is to adopt some stacking technique and investigate the observational signal statistically (e.g., Planck Collaboration et al. 2013d; Wang et al. 2014b; Ma et al. 2015; Lim et al. 2018a).

The baryonic gas in filaments and sheets is more diffuse and colder, and is thus even harder to investigate using X-ray emission and the tSZ effect. Significant detections of the tSZ signal and X-ray emission have been obtained only for massive



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filaments between pairs of massive merging clusters (e.g., Planck Collaboration et al. 2013b; Sugawara et al. 2017; Reiprich et al. 2021). Faint X-ray filaments around individual clusters are detected for a few clusters with extremely deep observations (Eckert et al. 2015; Connor et al. 2018). By stacking a large number of galaxy pairs, one may measure the weak tSZ effect produced by the gas between galaxies (e.g., de Graaff et al. 2019). However, the interpretation of the stacking results is not straightforward, because it is unclear whether the signal is dominated by the diffuse gas associated with the filaments or by small halos embedded in the filaments.

The gas components can also be observed through the absorption lines that they produce in the spectra of background sources like quasars, such as Ly α , Mg II, and O VI absorption line systems (e.g., Penton et al. 2000; Tumlinson et al. 2011b; Rudie et al. 2012; Shull et al. 2012; Werk et al. 2013; Danforth et al. 2016). Some of the observed absorption lines are clearly associated with galaxies, indicating that they are caused by the gas around galaxies (gaseous halos) or even by the gas associated with outflows driven by stellar and AGN feedback (e.g., Rudie et al. 2013; Lan & Mo 2018). Filament gas that is not physically connected to any halos can also produce observable absorption line systems (e.g., Nicastro et al. 2018; Pessa et al. 2018). Other methods, such as emission lines and radio emission, have also been used to study gas in galaxy groups and clusters (see, e.g., Brown & Rudnick 2011; Zhang et al. 2016; Gheller & Vazza 2020). Despite all these, a large fraction of the gas, about 30% to 40% of the baryons, remains undetected in the low-redshift universe (e.g., Bregman 2007; Shull et al. 2012), and the details of the physical processes responsible for the evolution and state of the gas are still unclear.

Theoretically, numerical simulations play a critical role in understanding the physical processes that determine gas properties in the cosmic web. Hydrodynamic simulations show that the gas in the cosmic web is processed and heated by strong shocks during the formation of halos and filaments (Kereš et al. 2005; Dekel & Birnboim 2006; Kang et al. 2007; Bykov et al. 2008; Schaal & Springel 2015; Zinger et al. 2018). In particular, more than 40% of the baryonic gas at $z \sim 0$ is found to be located outside halos, in a warm-hot intergalactic medium (WHIM) with a temperature between 10^5 and 10^7 K (e.g., Cen & Ostriker 1999; Davé et al. 2001; Haider et al. 2016; Cui et al. 2019; Martizzi et al. 2019). Most of the gas in the WHIM tends to reside in filaments, and the relatively low density and high temperature makes the gas hard to detect in both emission and absorption. Simulations also show that AGN feedback, stellar winds, and ram pressure stripping can not only affect the thermodynamic properties and spatial distribution of the gas in the cosmic web (see, e.g., Sun et al. 2009; Lovisari et al. 2015; Lim et al. 2018b; Amodeo et al. 2021; Schaan et al. 2021; Truong et al. 2021) but also chemically enrich the circumgalactic medium (CGM) and intergalactic medium (IGM; e.g., Cen & Ostriker 2006; Liang et al. 2016; Oppenheimer et al. 2016; Rahmati et al. 2016; Nelson et al. 2018; Boselli et al. 2022).

Clearly, to constrain models using observations, it is necessary to compare model predictions and observational data in a meaningful way. In general, the comparison is made in a statistical sense, in terms of summary statistics, such as distribution functions and correlations. Since both observational and simulation samples are finite, one needs to deal with the cosmic variance to achieve an unbiased comparison

between models and data. This is a serious challenge for both observations and simulations. For example, large-scale structures in the cosmic web, such as filaments and massive clusters are rare, so that the observational samples are usually small; observations of absorption systems are limited by the background sources available, so that only the gas distribution along a limited number of lines of sight can be sampled. All these make it difficult to obtain a statistically fair sample for the objects concerned. Hydrodynamic simulations are limited by the dynamic range they cover, and the simulation volume is usually relatively small, making it difficult to sample fairly the large-scale structure of the cosmic web. It is thus imperative to have as much theoretical and empirical input as possible to help compare the model predictions with the limited amount of observational data in an unbiased way. The bias can be minimized if comparisons are made for systems that have both the same environments and the same formation histories. This can be achieved if one can accurately reconstruct the initial conditions for the formation of the structures from which the actual gas emission and absorption are produced. Such a reconstruction is now possible (e.g., Hoffman & Ribak 1991; Nusser & Dekel 1992; Frisch et al. 2002; Klypin et al. 2003; Jasche & Wandelt 2013; Kitaura 2013; Wang et al. 2014a; Sorce et al. 2016; Bos et al. 2019; Horowitz et al. 2019; Modi et al. 2019; Kitaura et al. 2021). As shown in Wang et al. (2016), constrained simulations (CSs) using such reconstructed initial conditions can provide reliable representations of the structure, motion, and formation history of the objects observed in the real universe. Hydrodynamic simulations can also be carried out with these reconstructed initial conditions to follow the evolution of the gas component, facilitating unbiased comparisons between model predictions and observational data. CSs can also provide reliable tracers of the mass distribution in the observed universe, which is important for stacking analyses to discover unidentified gas in the cosmic web (e.g., Lim et al. 2018b, 2020).

In this paper, we use hydrodynamic simulations to follow the evolution of the gas component in three volumes covered by the Sloan Digital Sky Survey (SDSS; York et al. 2000), using initial conditions reconstructed from the ELUCID project (Wang et al. 2014a, 2016). The three volumes are chosen to contain the Coma cluster, SDSS Great Wall, and a low-density *void*. The main goal of this paper is to demonstrate the power and uniqueness of CSs in investigating and understanding the gas component in the observed cosmic web. The paper is organized as follows. Section 2 describes the three constrained hydrodynamic simulations and the method to characterize the cosmic web. Section 3 shows the evolution of gas properties in and around these large-scale structures in the local universe. We analyze the properties of the gas and study how they are correlated with the local properties of the cosmic web in Section 4. We discuss how to use CSs to facilitate unbiased comparisons between model predictions and observational data in Section 5. Finally, we summarize and discuss our results in Section 6.

2. Constrained Simulations of the Local Universe

2.1. Initial Conditions

To simulate the evolution of the large-scale structure observed in the local universe, we use initial conditions reconstructed from the ELUCID project in the volume covered

Table 1
The Properties of the Three Simulated HIR

Simulation	α_J	δ_J	z	Geometry	Size	ρ_m	N_g	N_{d1}	N_{d2}	N_{d3}
CM	194.8	27.9	0.0241	sphere	$R = 30 h^{-1}\text{Mpc}$	1.39	7.04e8	7.04e8	6.17e8	6.18e8
GW	188.0	2.0	0.0775	cuboid	$75 \times 30 \times 30 (h^{-1}\text{Mpc})^3$	2.19	9.91e8	9.91e8	8.67e8	8.66e8
LD	222.6	45.5	0.0519	sphere	$R = 40 h^{-1}\text{Mpc}$	0.57	1.09e9	1.09e9	9.56e8	9.56e8

Note. The table contains the location of the HIR center (α_J , δ_J , and redshift z), geometry, comoving size, and mean density (ρ_m , in units of the cosmic mean density). N_g and N_{d1} are the numbers of gas and dark matter particles in the iHIR; N_{d2} and N_{d3} are the numbers of dark matter particles in the LIR1 and LIR2 regions, respectively. The sky coverage of the three CSs is shown in Figure 1.

by the SDSS galaxy redshift survey. The reconstruction consisted of four steps. First, we identified galaxy groups from the SDSS galaxy catalog (Yang et al. 2005, 2007). Second, we corrected the distance of each group for redshift distortions using linear perturbation theory (Wang et al. 2012). Third, we reconstructed the present-day mass density field using a halo-domain method (Wang et al. 2009, 2016). Finally, we employed an HMC+PM technique to recover the initial conditions from the present-day density field (Wang et al. 2013, 2014a). Wang et al. (2016) performed a large N -body simulation using these reconstructed initial conditions in the SDSS Dr7 (Abazajian et al. 2009) survey volume at $z \leq 0.12$. In what follows, we refer to this CS as the original CS (hereafter OCS). A comprehensive description of these steps and the OCS can be found in Wang et al. (2016), and we refer readers to that paper for details.

In this paper, we focus on three regions that contain special large-scale structures of interest. We adopt a zoom-in technique, in which the interested structures are simulated with a higher resolution than other regions in the simulation volume (Katz & White 1993). To do this, we select, for each zoom-in simulation, a high-resolution region (HIR) to contain the particular structure of interest from the $z = 0$ snapshot of the OCS. The HIR is either a cuboid or a spherical volume (see below and Table 1). We follow all the particles in the HIR back to an initial time. We then select an initial HIR (hereafter iHIR) that has exactly the same comoving center, shape, and orientation (for a cuboid geometry) as the HIR. The iHIR is set to contain all particles in the HIR, and so the comoving size of the iHIR is larger than that of the HIR. To prevent lower-resolution particles from entering the HIR region in the subsequent evolution, a buffer region is used in defining the iHIR. The size of the buffer region is set to be 5% of the iHIR, so that the iHIR volume is enlarged by 15.8%.

The iHIR is sampled with high-resolution particles, and the cosmic density field outside the iHIR is sampled with particles of three successively lower resolutions. Except for the lowest-resolution region, the other two low-resolution regions have exactly the same geometry and center as the iHIR. The size of the first low-resolution region (hereafter LIR1) is twice that of the iHIR, and its mass resolution is 8 times lower so that the number of LIR1 particles is similar to that in the iHIR. The size and mass resolution of the second low-resolution region (LIR2) are chosen to be twice as large and 8 times as low as those of the LIR1, respectively. Finally, the rest of the simulation volume is sampled with particles, each with a mass that is 8 times as large as that of a LIR2 particle. To set up the zoom-in initial condition, we generate a displacement field for high-resolution particles with a grid of 6144^3 cells in the whole simulation box. In the three lower-resolution regions, we bin the high-resolution particles according to the corresponding

mass resolution. Only HIRs contain gas particles. We thus split each high-resolution particle into one DM particle and one gas particle. The mass ratio of the two particles equals to $(\Omega_{m,0} - \Omega_{b,0})/\Omega_{b,0}$, and their separation is half of the mean separation of particles. Our inspection shows that this choice of resolution hierarchy ensures that no lower-resolution particle enters the HIR region. For reference, we list the particle numbers used in iHIR, LIR1, and LIR2 in Table 1.

The three HIRs that we choose to simulate are the following. The first one contains a part of the SDSS Great Wall at redshift ~ 0.08 . Since the Great Wall is a long filamentary structure, we choose a cuboid HIR with size $75 \times 30 \times 30 (h^{-1}\text{Mpc})^3$. Hereafter, we refer to this zoom simulation as the GW simulation. The second one is centered on the Coma galaxy cluster (CM) at a redshift of 0.0241. The HIR is set to be a spherical volume with a radius of $30 \text{ Mpc } h^{-1}$. We refer to the corresponding zoom simulation as the CM simulation. The third one is a low-density region at $z \sim 0.05$, and is chosen to be a spherical volume with a radius of $40 \text{ Mpc } h^{-1}$. We refer to this simulation as the LD simulation. Figure 1 shows the location of the three selected regions in the J2000.0-coordinate system represented by R.A. (α_J) and decl. (δ_J). The mean mass densities in the three HIRs are 2.19 (GW), 1.39 (CM), and 0.57 (LD) times the cosmic mean density. In Table 1, we list the basic information of the three HIRs.

We adopt cosmological parameters from WMAP5 (Dunkley et al. 2009), the same as the OCS: $\Omega_{\Lambda,0} = 0.742$, $\Omega_{m,0} = 0.258$, $\Omega_{b,0} = 0.044$, $h = H_0/100 \text{ km s}^{-1} \text{ Mpc} = 0.72$, $\sigma_8 = 0.80$. The mass resolution of the HIRs is 8 times higher than the OCS simulation. The corresponding masses of DM and gas particles in the HIR are $3.20 \times 10^7 h^{-1} M_\odot$ and $6.74 \times 10^6 h^{-1} M_\odot$, respectively. The mass resolution of our HIRs is close to that used in Huang et al. (2020). The masses of DM particles in LIR1, LIR2, and the rest part of the simulation are 3.09×10^8 , 2.47×10^9 , and $1.98 \times 10^{10} h^{-1} M_\odot$, respectively. All three simulations follow the evolution of the cosmological density field in a periodic box of comoving length $500 h^{-1} \text{ Mpc}$, and the initial redshift of each simulation is set at $z = 120$. Softening lengths for high-resolution particles are all set to be 1.8 comoving $h^{-1} \text{ kpc}$.

2.2. Simulation Code

We run all the simulations using Gadget-3, an updated version of Gadget-2 (Springel 2005), as described in Huang et al. (2019, 2020). The gravitational forces are evaluated using a particle mesh and oct-tree algorithm. The code includes several recent numerical improvements in the smoothed particle hydrodynamics (SPH) technique (Huang et al. 2019). To summarize, we use the pressure-entropy formulation (Hopkins 2013) of SPH to integrate the fluid equations and a quintic spline kernel to measure the fluid quantities over 128

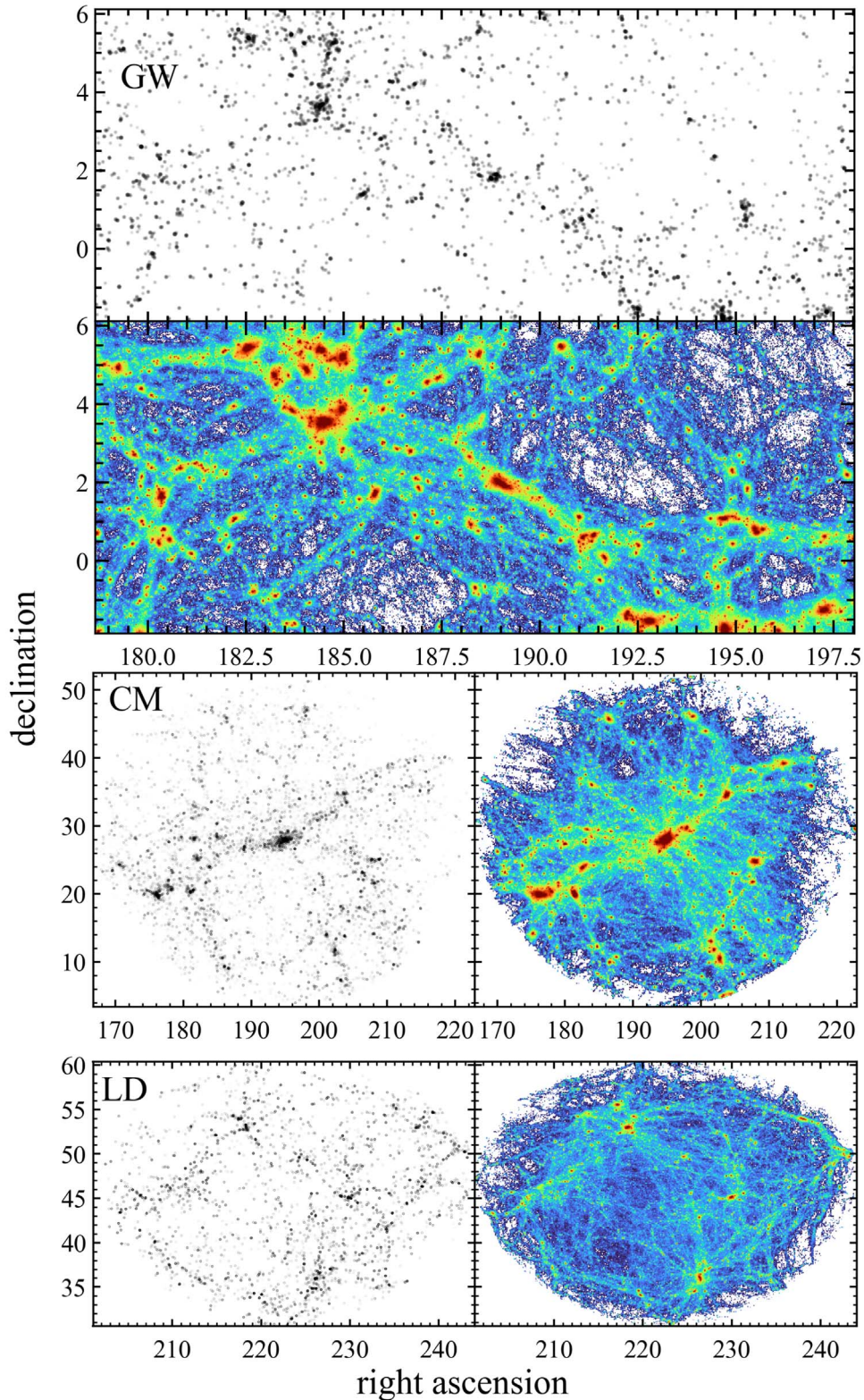


Figure 1. These six panels show the spatial distribution of galaxies (black dots) and dark matter (color-coded maps) in the constrained simulations of the Great Wall region (GW, top two panels), the Coma cluster region (CM, middle two panels), and the low-density void region (LD, bottom two panels) in the J2000.0 R.A. and decl.. Note that the upper panel of GW and the left panels of CM and LD show real galaxies from the SDSS galaxy redshift survey.

neighboring particles. We also use the Cullen & Dehnen (2010) viscosity algorithm and artificial conduction as in Read & Hayfield (2012) to capture shocks more accurately and to reduce numerical noise. Both the artificial viscosity and the conduction are turned on only in converging flows with

$\nabla \cdot \mathbf{v} < 0$ to minimize unwanted numerical dissipation. We also include the Hubble flow while calculating the velocity divergence. Our fiducial code leads to considerable improvements in resolving the instabilities at fluid interfaces in subsonic flows and produces consistent results with other

state-of-the-art hydrodynamic codes in various numerical tests (Sembolini et al. 2016a, 2016b; Huang et al. 2019).

Besides including cooling from hydrogen and helium and the effects of a photoionizing UV background field (Haardt & Madau 2012), we also add metal line cooling including photoionization effects for 11 elements as in Wiersma et al. (2009), and we recalculate the cooling rates according to the ionizing background (Haardt & Madau 2012). The star formation processes are modeled as in Springel & Hernquist (2003), which includes a subgrid model for the multiphase interstellar medium (ISM) in dense regions with $n_{\text{H}} > 0.13 \text{ cm}^{-3}$, and a star formation recipe that is scaled to match the Kennicutt–Schmidt relation. In this paper, we will distinguish SPH particles as galaxy particles based on whether or not their densities are higher than this density threshold. We specifically trace the enrichment of four metal species C, O, Si, Fe that are produced from type II SNe, type Ia SNe, and AGB stars as in Oppenheimer & Davé (2008). These processes also generate energy that we add to the simulations as thermal energy. However, the input energy from these feedback processes only has subdominant effects on galaxy formation compared to the wind feedback (Oppenheimer & Davé 2008), as it is typically quickly radiated away.

As in Huang et al. (2019, 2020), we adopt a kinetic subgrid model for the stellar feedback. We implemented the new wind algorithm into our SPH code based on GADGET-3 (see Springel 2005 for reference). SPH particles in star-forming regions have some probability of being ejected from their host galaxy in the direction $\mathbf{v} \times \mathbf{a}$, implemented with a one-at-a-time particle ejection algorithm. We temporarily shut off the hydrodynamic forces to allow particles to escape the dense ISM. While this is artificial, it simply recognizes the fact that feedback *does* drive outflows in real galaxies (and in ultra-high-resolution simulations of isolated galaxies, like those of Hopkins et al. 2012), rather than being locally thermalized and radiated away, owing to collective effects in a multiphase ISM.

We adopt a variant of the wind algorithm implemented in GADGET, using momentum-driven wind models very similar to those found in high-resolution zoom simulations. Building on the work of Oppenheimer & Davé (2006), we model the launch of galactic winds from star-forming galaxies with controlled parameters of the mass loading factor, $\eta \equiv \text{ejection rate/SFR}$ ($\eta \sim \sigma_{\text{gal}}^{-\beta}$ for small σ_{gal} and $\eta \sim \sigma_{\text{gal}}^{-\alpha}$ for large σ_{gal}) and the initial wind speed, v_w , that scales linearly with the velocity dispersion of the galaxy (σ_{gal}). We identify galaxies on the fly at intermittent time steps during a simulation using a friend-of-friend (FoF) group finder, which at the same time computes the properties of these galaxies and estimates the dispersion using the total mass of the galaxy.

We adopt the same set of wind parameters as the fiducial simulation from Huang et al. (2020), which was the most successful wind launch scalings, in terms of matching a broad range of observations (e.g., Davé et al. 2013), where $\alpha = 1$ and $\beta = 3$, i.e., \sim momentum-driven wind scalings for large σ_{gal} and supernova-energy-driven wind scalings for small σ_{gal} , and $V_w \sim \sigma_{\text{gal}}$. These scalings are very similar to those found in very-high-resolution galaxy zoom simulations (Hopkins et al. 2012, 2014; Muratov et al. 2015). Our wind scalings are at wind launch from the star-forming regions of the galaxy while the very-high-resolution zoom simulations report their wind scalings at $0.25r_{\text{vir}}$ (Muratov et al. 2015). Hence, we have had

to slightly increase our wind launch velocities to reproduce their behavior at $0.25r_{\text{vir}}$. This seemingly minor change has important effects; if we were to simply apply the Muratov et al. (2015) scalings at launch, as in all past work, then owing to gravitational deceleration many wind particles in moderately large galaxies would not even reach $0.25r_{\text{vir}}$, and those that do would have a V_w that is almost independent of σ_{gal} . We also cap the wind speed so that the energy in the winds does not exceed that available in the supernova.

Feedback from AGNs is not included in our simulations; we will discuss the potential impact of not including AGNs later. Since we are mainly interested in gas properties at large scales, the impact of AGN feedback may be relatively small.

2.3. Characterizing the Cosmic Web

Halos are identified using a FoF algorithm (Davis et al. 1985) with a linking length b equal to 0.2 times the mean DM particle separation. The FoF algorithm is first applied to high-resolution DM particles to identify DM halos. A gas or star particle is assigned to the same halo defined by the DM particles if its distance to the nearest DM particle is less than b . Halos with at least 20 DM particles are identified, and the halo mass, M_h , is defined as the sum of the masses of all particles in the halo. Galaxies, including star and star-forming gas particles, are identified using the Spline Kernel Interpolative Denmax algorithm (Kereš et al. 2005).

In addition to halos, we are also interested in other components of the cosmic web. We adopt the so-called ‘‘T-Web’’ method (e.g., Hahn et al. 2007) to classify the cosmic web. The method is based on the eigenvalues of the tidal tensor defined as

$$\mathcal{T}_{ij} = \partial_i \partial_j \phi, \quad (1)$$

where ϕ is the peculiar gravitational potential, and it obeys a modified Poisson equation:

$$\nabla^2 \phi = \delta. \quad (2)$$

Note that this equation is scaled by $4\pi G\bar{\rho}$, and δ is the mass overdensity defined over grid cells and smoothed on a comoving scale of R_s . The three eigenvalues of \mathcal{T}_{ij} are denoted by λ_i , where $i = 1, 2, \text{ and } 3$, with $\lambda_1 \geq \lambda_2 \geq \lambda_3$. Following common practice, we define knots as locations where all three eigenvalues are above a given threshold value λ_{th} , while filaments, sheets, and voids are locations where two, one, and none of the eigenvalues are above λ_{th} , respectively.

We follow Hahn et al. (2007) and adopt $\lambda_{\text{th}} = 0$. Since we want to classify the cosmic web on small scales, we choose a smoothing scale of $R_s = 0.5 h^{-1} \text{Mpc}$, much smaller than the $2.0 h^{-1} \text{Mpc}$ often adopted in previous studies. We will discuss the impact of using different smoothing scales. Thus, the knots defined here include not only massive galaxy clusters but also small halos; only the most underdense regions are classified as voids, as we will see below.

To distinguish gas in different components of the cosmic web, we refer to the gas in halos as halo gas, and the gas outside halos as diffuse gas. The diffuse gas is further separated into knot gas, filament gas, sheet gas, and void gas. In the following analysis, we exclude gas particles within galaxies, defined as gas that has $n > 0.13 \text{ cm}^{-3}$ and a temperature less than $10^{4.5} \text{K}$, as we are only interested in gas properties on larger scales.

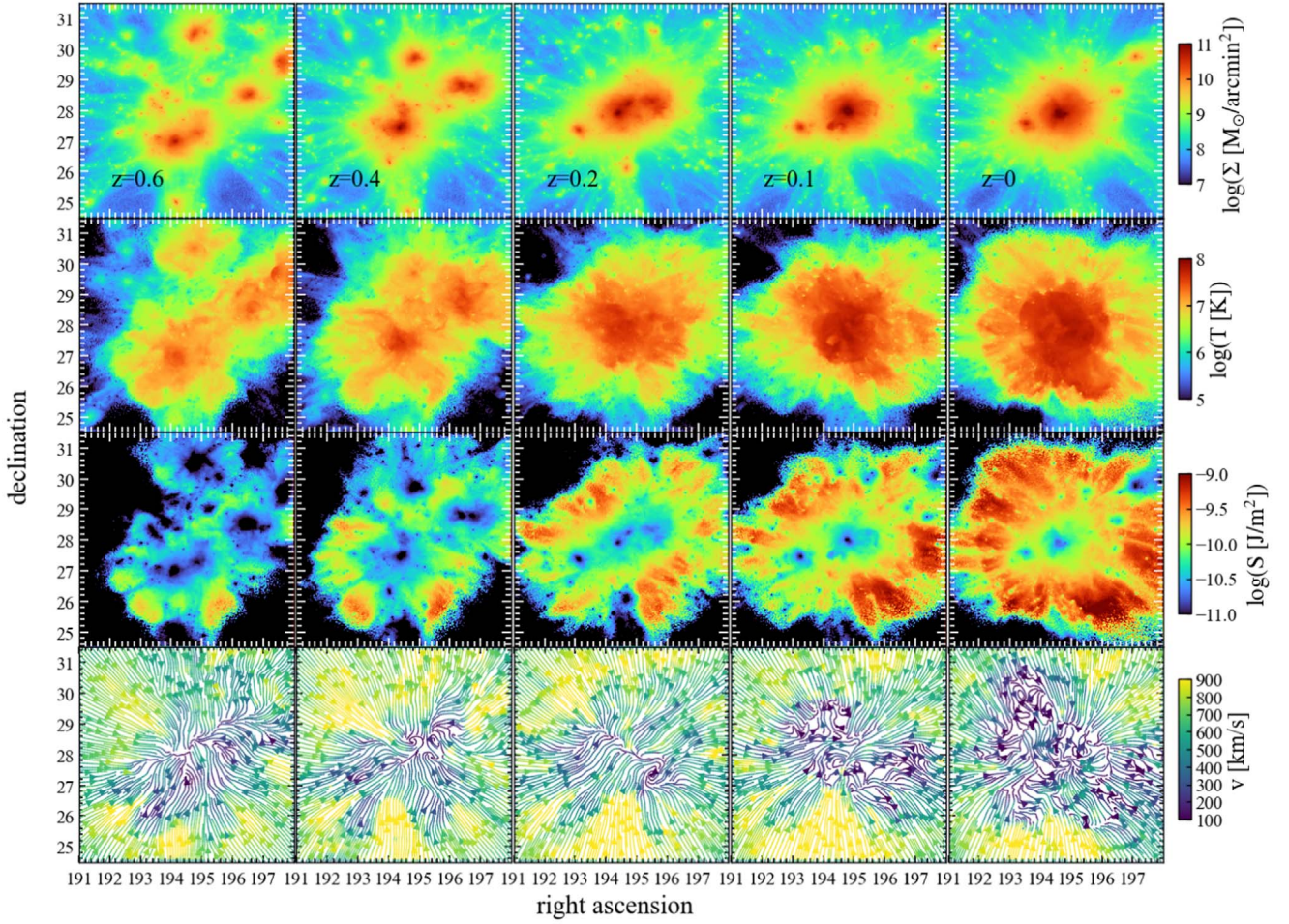


Figure 2. The evolution of the Coma galaxy cluster in the CM simulation from $z = 0.6$ to $z = 0$. The top shows the surface mass density of gas particles, the second row shows the gas temperature, the third row shows the gas entropy, and the bottom row shows the velocity component that is perpendicular to the line of sight. Note that the mean velocity is removed.

3. Evolution of Large-scale Structures in the Local Universe

Figure 1 shows the spatial distribution of SDSS galaxies in the three selected regions. Galaxies are viewed from Earth and shown in the J2000.0-coordinate system represented by the R.A. (α_j) and decl. (δ_j). The present-day density fields of the corresponding CSs are also plotted for comparison. As one can see, most of the massive clusters and filaments and large voids are well reproduced in the CSs, as well as some small filaments. We refer readers to Wang et al. (2016) for a detailed discussion about the statistics of the OCS and its comparison with the distributions of real galaxies. In what follows, we focus on the spatial distribution of gas properties, such as mass, temperature, entropy ($s \equiv kT/n^{2/3}$), and velocity, as well as the time evolution of these properties.

3.1. The Coma Galaxy Cluster (CM) Simulation

We first show the evolution of the gas in a region around the CM (Figure 2). As one can see, the Coma cluster has experienced a number of violent major mergers during the past 6 billion years. In particular, the most recent major merger occurred at $z \sim 0.2$. Most of the major mergers occurred along filaments connecting the Coma cluster with other galaxy clusters as shown in Figure 1. This is consistent with the fact that the Coma cluster has two bright dominant galaxies, likely

the central galaxies of the two massive progenitors of the last major merger (at $z \sim 0.2$). Our simulation also reveals an ongoing minor merger event on the left side of the cluster ($\alpha_j \sim 193.5$ and $\delta_j \sim 27.5$). This is similar to the observed galaxy distribution represented by the subgroup, NGC 4839. At $z = 0$, the simulated Coma cluster has a mass $M_{200} = 7.52 \times 10^{14} h^{-1} M_\odot$ and a virial radius $r_{200} = 1.48 h^{-1} \text{Mpc}$, corresponding to $\sim 70'$, where M_{200} is defined as the mass contained in a spherical region of radius r_{200} , within which the mean mass density is 200 times the critical density. For comparison, the real Coma cluster has a radius of $r_{200} = 70'$ according to a scaling relation shown in Simionescu et al. (2013).

At $z = 0.2$ when the major merger starts, the velocity stream lines appear to converge at the center of the cluster, as shown in the bottom panels of Figure 2. One can see clearly that two clumps of gas are colliding with a relative velocity larger than about 1200 km s^{-1} . Such a collision is expected to produce two strong shocks propagating in opposite directions (see, e.g., Zinger et al. 2018). This is signified by the sharp temperature jump in the inner region seen in the temperature map at $z = 0.1$. The shock propagates outward and, at $z = 0$, reaches a scale larger than the virial radius of the Coma cluster. Meanwhile, a new bow-like shock is produced in the inner region (see the density map at $z = 0$), indicating that the cluster has not yet relaxed to hydrodynamic equilibrium. This result is consistent

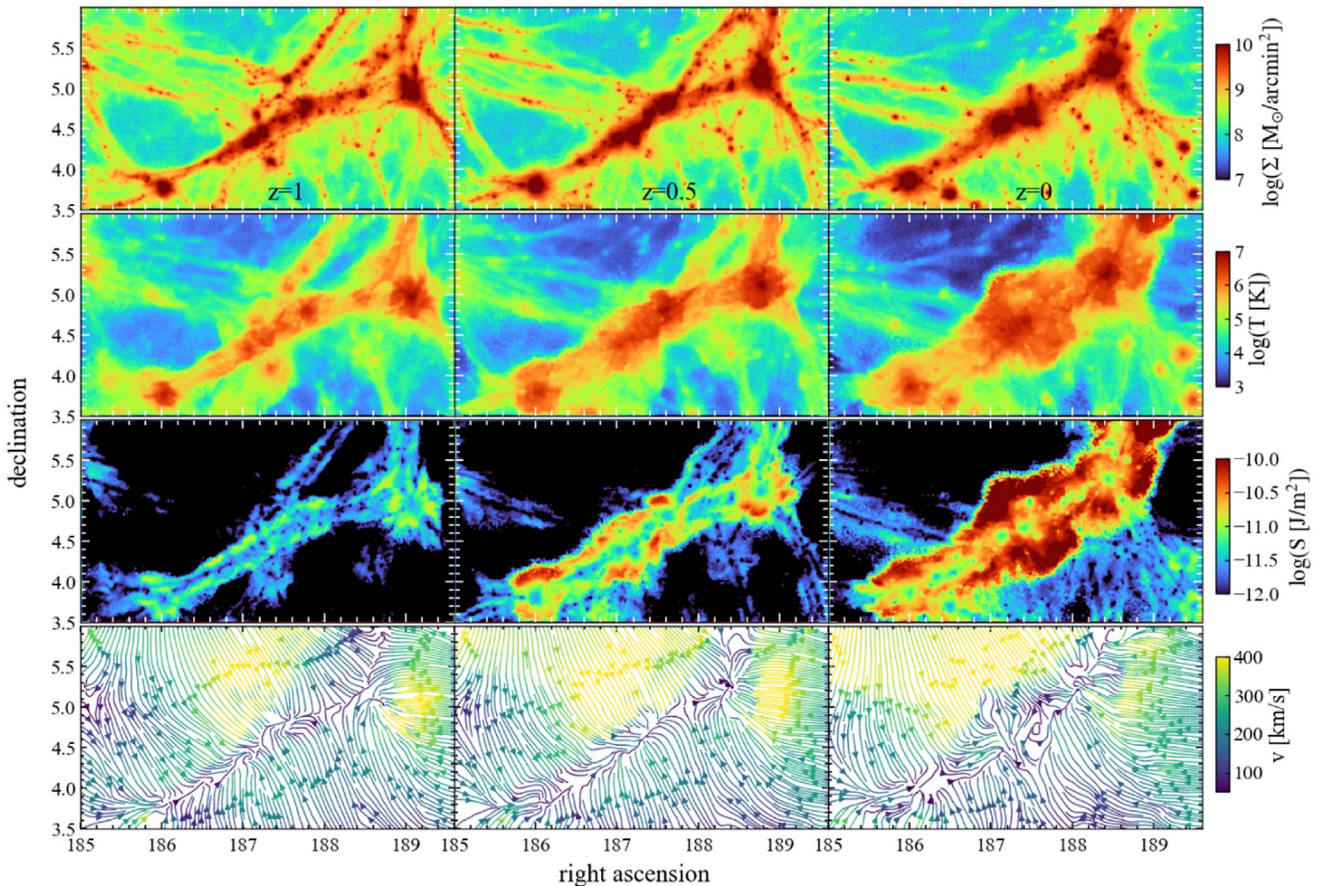


Figure 3. Similar to Figure 2 but for a massive filament in the GW simulation. Here we present the results from $z = 1$ to 0. Note that this filament is only a small part of the SDSS Great Wall.

with the results obtained from recent X-ray observations (e.g., Mirakhor & Walker 2020; Churazov et al. 2021), and we will come back to this later in Sections 5.1 and 5.2.

The outgoing shocks can be seen to collide subsequently with the gas accretion and heat the gas, as shown in the entropy map. This interaction makes the gas stream lines complex. As one can see from the $z = 0$ velocity map, the infall velocity along the filament (from the southwest to the northeast direction) is larger than that perpendicular to the filament. This indicates that gas is still being accreted along the filament, while, perpendicular to the filament, a larger fraction of the gas is driven by outward shocks.

3.2. The SDSS Great Wall (GW) Simulation

The gas evolution around massive filaments looks quite different from that around galaxy clusters. In Figure 3, we show the evolution of a massive filament in the GW simulation. Note that this filament is only a small part of the Great Wall. The whole filamentary structure can already be seen at $z > 1$. It is consistent with the results presented in Wang et al. (2016). The filament thickens itself by accreting gas from nearby underdense regions. The velocity streams around the filament have remained at a value of 300 to 400 km s^{-1} since $z = 1$ and are roughly perpendicular to the filament. The gas accretion onto the filament also leads to accretion shocks. As one can see, the gas entropy at the boundary of the filamentary structure is higher than that in the inner part of the filament, and the entropy difference increases with decreasing redshift,

indicating that accretion shocks constantly heat the gas surrounding the filament. A similar evolution can also be seen for the small filaments shown in the figure.

The structure and evolutionary history shown in Figure 3 indicate that the diffuse gas in filaments is generally heated by accretion shocks, and the heating effect becomes stronger as the structure grows. Since the velocity of the gas is lower than that around the Coma cluster, the shocks generated are weaker. For example, for the main filament shown in the figure, the temperature can reach 10^6 – 10^7 Kelvin. The velocity map shows that the gas flows in the filaments are quite turbulent, which may provide an important process to mix metal-rich gas generated by galaxies with the IGM.

3.3. The Low-density Region (LD) Simulation

In the center of the selected low-density region at $z \sim 0.05$, there is a large void structure. Figure 4 shows that the structure in the void grows very slowly, as expected from the low mass density that drives gravitational instability. Gas particles are seen to flow continuously out from the void region, making the density decrease with time over most of the volume. The gas temperature in the void also declines slowly with time. The gas entropy, which is roughly homogeneous over the void region, increases with time. This is consistent with the fact that gas in low-density regions is rarefied by the expansion of the universe and is heated by the UV background. As one can see in the density and temperature maps, abundant small filamentary structures can be seen since $z > 2$ and the gas temperature in

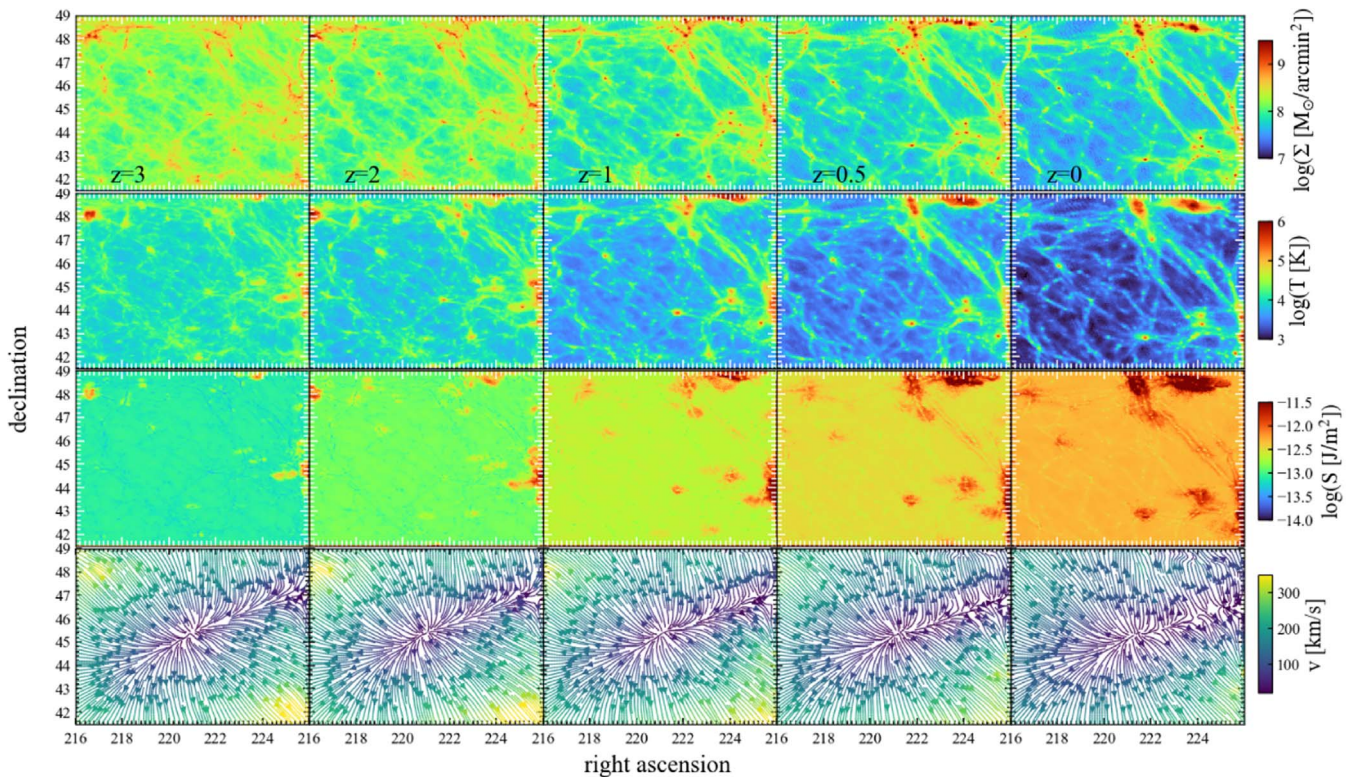


Figure 4. Similar to Figure 2 but for the central void region in the LD simulation. Here we present the results from $z = 3$ to 0.

them has remained in the range of $10^{3.5} - 10^{4.5}$ K. However, most of these filaments are barely visible in the entropy map, indicating that the formation of these filamentary structures is an adiabatic process in an ambient medium of similar specific entropy. From $z \sim 3$ to the present day, almost all the small filaments have an entropy slightly lower than their ambient media, indicating that some radiative cooling may be present within them. At $z \sim 0$, the entropy of some relatively large filaments appears to be higher than the environment, because shock heating starts to become significant. Clearly, filaments of different sizes have gone through different heating processes.

4. Gas Properties and Their Correlation with the Cosmic Web

The above analysis suggests that the gas in different components of the cosmic web may be heated by different processes. In Figure 5, we show the gas temperature distribution in different components of the cosmic web, as classified using the T-Web method described in Section 2.3. The distribution in a given component is defined as $dm_s/d \log T/m_{\text{tot}}$, where dm_s is the mass of gas particles with a temperature in the range of $\log T$ to $\log T + d \log T$ in the component, while m_{tot} is the total gas mass, rather than that in the component in question. Note that only gas particles in the HIRs are used for the analysis. Thus, the results shown can be used to understand not only the temperature distribution within a given type of component but also the contribution from different components.

The black lines in the three panels of Figure 5 show the results for all the gas in the three HIRs, respectively. The gas temperature distributions in the three regions are clearly different. In the GW simulation, there is a significant bump around $\log(T/K) \sim 7$, which is absent in the other two

regions. This is produced by the existence of the massive filamentary structure in the GW. In the CM simulation, the gas temperature can reach as high as 10^8 K because of the presence of the massive Coma cluster. Finally, the LD simulation is dominated by low-temperature gas, as expected.

The temperature distribution for gas outside halos (diffuse gas) is also shown in each of the three panels. As one can see, the spike at $\log(T/K) \sim 4$ for all gas particles disappears, indicating that this feature is dominated by cold CGM in halos. The diffuse gas dominates at the low-temperature end and becomes less important as the temperature increases. At $\log(T/K) = 5$, more than 80% of the gas is in the diffuse component, and the fraction decreases to about 20% at $\log(T/K) = 7$. About half of the gas with $5 < \log(T/K) < 7$ (the WHIM) is located outside halos, quite independent of the large-scale environment. This result is consistent with that obtained before (e.g., Cen & Ostriker 1999; Davé et al. 2001; Martizzi et al. 2019).

In Figure 5, we also plot the temperature distribution of the diffuse gas separated into the four different components of the cosmic web. In most cases, the gas in filaments dominates the diffuse gas at $\log(T/K) > 4$. This indicates that most of the warm-hot diffuse gas resides in filaments. At $\log(T/K) < 4$, sheet gas contributes the most. Void gas has a very narrow distribution peaked at the lowest temperature, and its contribution is important only in the LD simulation. Our results also show that the diffuse gas with the highest temperature is mostly associated with filaments rather than knots, which is different from the expectation that high-temperature gas is associated with the outskirts of massive clusters. We note that in detail the results depend on the smoothing scale R_s used for the analysis. If a larger value of R_s were used, the hot gas in the interior of halos could affect the estimate of gas properties in the outskirts

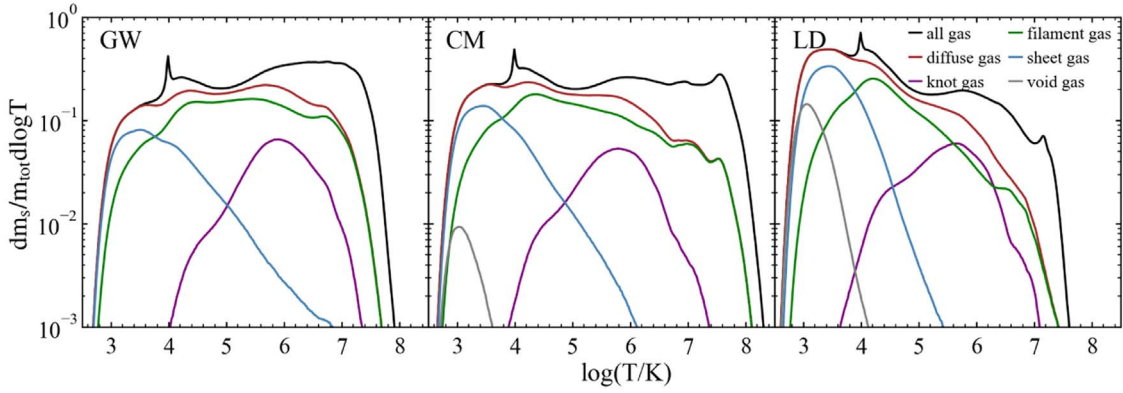


Figure 5. The temperature distribution of gas particles in different types of structures for the GW (left panel), CM (middle panel), and LD (right panel) simulations. Black lines are for all gas particles and dark-red lines for all diffuse particles. The other color-coded lines are for gas particles in different components of the cosmic web as indicated. See the text for details of the analysis.

of halos. For example, if we adopt $R_s = 2 h^{-1} \text{Mpc}$, instead of $R_s = 0.5 h^{-1} \text{Mpc}$ as used above, the diffuse gas associated with the knot component becomes hotter than the filament gas. The small smoothing scale, $R_s = 0.5 h^{-1} \text{Mpc}$, used here minimizes the contamination by halo gas in our analysis of the diffuse components.

Given the differences in the gas temperature between different components of the cosmic web as shown in Figure 5, it is interesting to directly check the correlation between the temperature and the local tidal field that was used to decompose the cosmic web. In Figure 6, we show the gas temperature as a function of $|\lambda_1|$ for halo gas and diffuse gas in the four components of the cosmic web.

As an example, here we only show the results for the GW simulation. The results for the other two simulations are similar, although there are differences in their temperature distributions (Figure 5).

As one can see, there is a strong and tight correlation between the gas temperature and the tidal field strength represented by $|\lambda_1|$. The correlation can be separated into three distinct components. For illustration, the two dashed lines show $\log T \propto 2 \log \lambda_1$ and $\log T \propto \log \lambda_1$, respectively. As one can see, the halo gas and the diffuse knot gas are both dominated by hot gas with $T \gtrsim 10^6 \text{K}$, and both follow the relation $\log T \propto \log \lambda_1$, suggesting that both are heated by similar processes. The gas in sheets is warm, with a temperature between 10^3 and 10^5K , follows $\log T \propto 2 \log \lambda_1$, and has almost no overlap with the halo and knot components. The gas in filaments has a more complex behavior. At high temperature, it follows the relation of the hot gas, while at low temperature, it obeys the correlation of the warm (sheet) component. The dividing point is around $\log T \sim 5.4$ where $\log |\lambda_1| \sim 1$. Finally, void gas, which is relatively cold, has a temperature quite independent of the tidal field strength.

The tight correlation shown above suggests that the local tidal field strength is a good indicator of the gas temperature before radiative cooling becomes important. It may thus be possible to reconstruct the 3D gas pressure distribution from the reconstructed total mass density field. To do this, we need to check how well the gas density correlates with the total mass density. Figure 7 shows the gas density versus mass density obtained by using a smoothing scale of $R_s = 0.5 h^{-1} \text{Mpc}$. The two densities are tightly and linearly correlated. To predict the pressure at a given location, we first use the value of λ_1 , again

smoothed with $R_s = 0.5 h^{-1} \text{Mpc}$, to predict the gas temperature using the median relation shown in Figure 6. We obtain the gas density, ρ_g , from the total mass density, ρ_m , using a simple relation $\rho_g = f_b \rho_m$, where f_b is the universal baryon fraction. The electron pressure at the grid point is obtained using $P_{e,\text{pre}} = kn_e T$, assuming fully ionized gas and primordial metallicity. In Figure 8, we show the comparison of the electron pressure predicted in this way with that obtained directly from the simulation. As one can see, the predicted pressure is tightly correlated with the simulated value, with a typical scatter of less than 0.5 dex, and the scatter increases with decreasing pressure. The method recovers the simulation results with only a small bias (up to 0.5 dex) that also increases with decreasing pressure. This dependence of the bias and scatter on pressure is caused by the weak correlation between tidal field strength and temperature at the low end of $|\lambda_1|$. For a typical filament with $\rho_g = 10 \bar{\rho}_b$ and $\log(T/\text{K}) = 5.5$, the electron pressure is about 0.71K cm^{-3} , and the scatter between the reconstruction and simulation is expected to be less than 0.25 dex. Thus, our pressure reconstruction is reliable at least for filaments and large sheets. As discussed in Section 5.1, this method could be applied to constrained, pure DM simulations to reconstruct the temperature, gas density, and pressure fields in regions where accurate reconstructions of the DM density field are available.

5. Connections to Observational Data

In this section, we use our CSs to generate a number of observable quantities, such as the expected thermal Sunyaev–Zel’dovich (tSZ) effect, X-ray emission, and HI and OVI absorption systems. These results can be compared with observational data statistically or for individual objects, taking full advantage of our CSs.

5.1. The Thermal SZ Effect

The tSZ effect provides an important avenue to investigate the ICM and IGM (Sunyaev & Zel’dovich 1972). The strength of the tSZ effect in a given direction is described by the dimensionless Compton parameter, y , defined as

$$y = \frac{\sigma_T}{m_e c^2} \int P(l) dl, \quad (3)$$

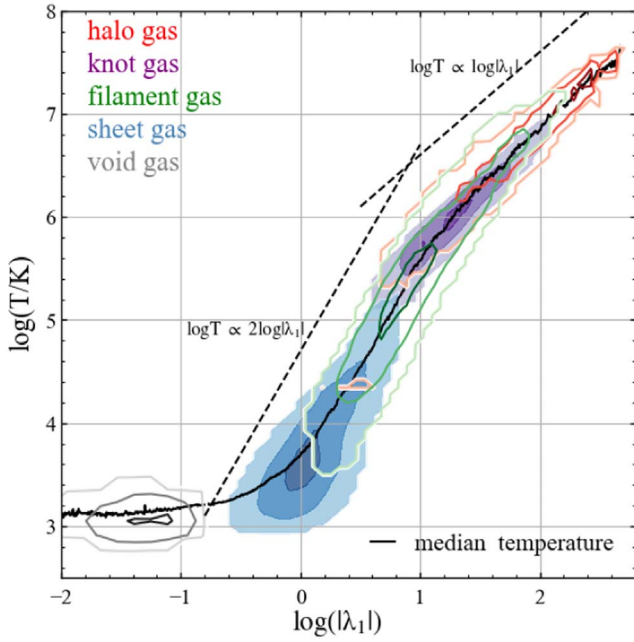


Figure 6. Gas temperature as a function of tidal field strength, $|\lambda_1|$, for different components of the cosmic web (and halos) at $z = 0$. Both temperature and tidal field are evaluated on grids. For each component, the contours enclose 10%, 45%, 80% of all the gas mass, respectively. Note that the results for individual components of the cosmic web do not include gas in halos, and that the halo gas shown here does not include gas particles in galaxies. The two dashed lines show two power-law functions with different slopes, as indicated, and the solid line shows the median temperature as a function of the tidal strength for all the gas particles.

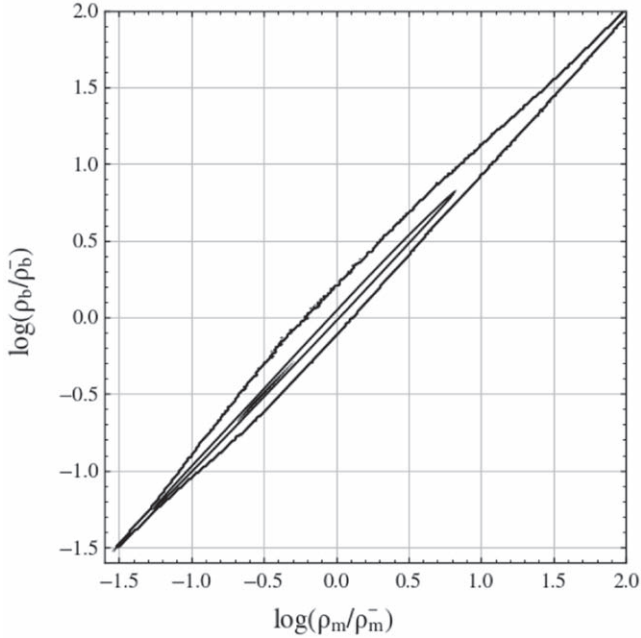


Figure 7. Mass density (ρ_m) vs. baryon density (ρ_b). Both densities are calculated in cubic cells of $(0.1 h^{-1} \text{Mpc})^3$ with a smoothing scale of $0.5 h^{-1} \text{Mpc}$, using the GW simulation, and scaled with their mean values, respectively. The contours enclose 10%, 80%, 99% of all cells.

where m_e is the electron mass, σ_T is the Thomson scatter cross section, c is the speed of light, P is the electron pressure, and the integration is along a given line of sight. For each simulation, we determine the ranges of α_J and δ_J , over which

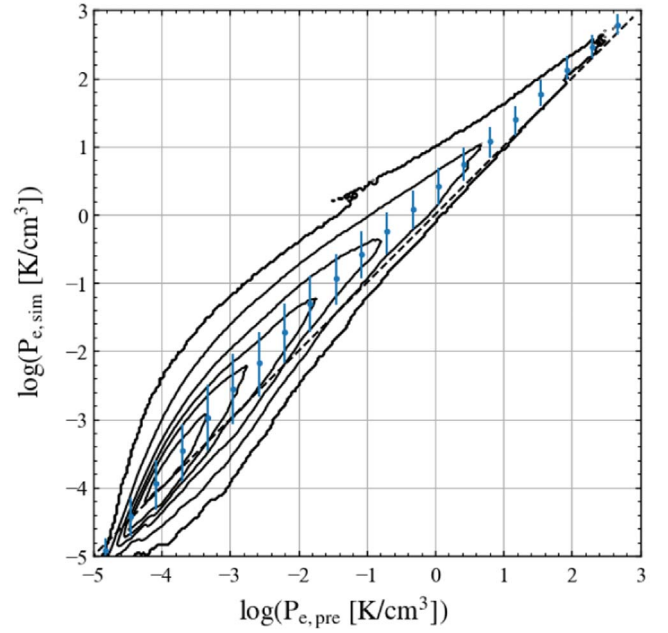


Figure 8. The predicted electron pressure ($P_{e,\text{pre}}$) vs. the simulated electron pressure ($P_{e,\text{sim}}$). The contours enclose, respectively, 10%, 30%, 50%, 70%, 90%, 99% of all the cells, each being a cube of $(0.1 h^{-1} \text{Mpc})^3$. The points show the median $P_{e,\text{sim}}$ as a function of $P_{e,\text{pre}}$, and error bars show the 16 and 84 percentiles. The dashed line corresponds to the one-to-one relation between the predicted and simulated electron pressures. Note that a typical filament with $\rho_b = 10\bar{\rho}_b$ and $\log(T/K) = 5.5$ has a pressure of 0.71K cm^{-3} .

the tSZ effect is to be modeled, and divide the sky into 2D pixels of a given angular size. We thus obtain a y -parameter map sampled on all the pixels in the sky coverage. This map is also used to study the profiles around individual objects, such as the CM.

We adopt two methods to construct y -maps. In the first one, we divide the volume covered by a 2D pixel into many small cells. The cell sizes are chosen to be smaller than the gravitational softening length of $1.8 h^{-1} \text{kpc}$. We estimate the number density and temperature of each cell from nearby gas particles using the SPH smoothing kernel. We then integrate along the line of sight (Equation (3)) to obtain the y -parameter of each pixel. The results presented in Figures 9 and 10 are obtained with this method (hereafter SphM). The SphM method is very time consuming and also unnecessary for some of our analyses. We therefore adopt a second, faster method for some of our presentations. Specifically, for a 2D pixel with a solid angle of $\Delta\Omega$ covering M gas particles, we approximate Equation (3) with

$$y = \frac{\sigma_T k_B}{m_e c^2 \Delta\Omega} \sum_i^M \frac{N_{e,i} T_i}{r_i^2}, \quad (4)$$

where $N_{e,i}$ and T_i are, respectively, the number of free electrons and the gas temperature represented by particle i , r_i is the distance of the particle, and k_B is the Boltzmann constant. This approach (hereafter ParM) is valid when the pixel size is much larger than the local SPH kernel radius. As shown in the Appendix, the ParM method is less accurate than the SphM, especially in low-density regions where the SPH kernel is large, but is sufficient for the purpose of this paper.

The left panel of Figure 9 shows the radial profile of the y -parameter of the Coma cluster, obtained from the simulated

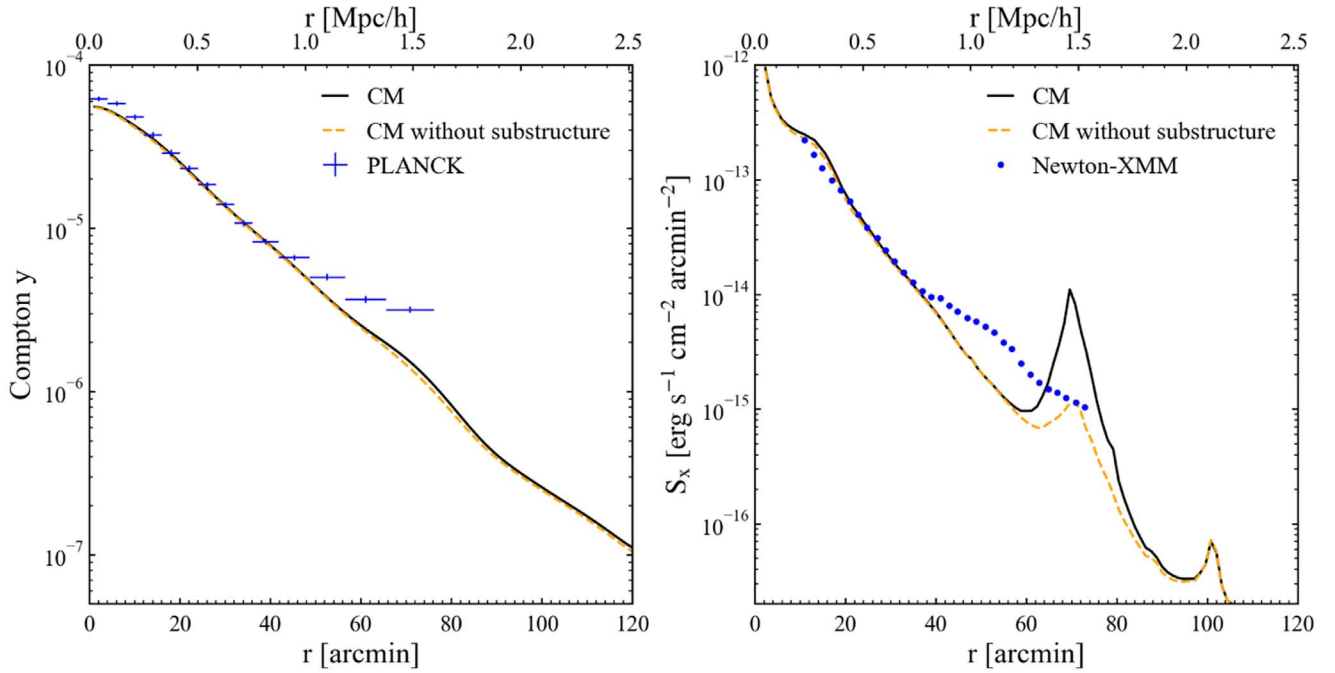


Figure 9. Left panel: the Compton y -profile (azimuthally averaged) for the Coma cluster. The data points, taken from Mirakhor & Walker (2020, private communication), are the measurements from Planck, and the lines show the results from our CM simulation. The solid black line shows the result averaged over the whole Coma cluster, while the dashed curve shows the result after the removal of the most significant substructure (see Figure 10). To compare with the observational data, the simulation result is smoothed with a FWHM = $10'$. Right panel: the surface brightness profile of X-rays in the band of 0.7–7.0 keV for the Coma cluster. The data points are based on Newton-XMM observations (Mirakhor & Walker 2020, private communication), and the two lines show results with and without the substructure removed, respectively.

y -map using SphM. The result is extended to an angular radius of $120'$, which is about 1.7 times the virial radius of the cluster. For comparison, we also show the observed y -parameter profile for the Coma cluster obtained in Mirakhor & Walker (2020) from the Planck y -map (e.g., Planck Collaboration et al. 2013c). Since the effective spatial resolution of the Planck y -map is about $10'$ FWHM, we smooth the simulation result with the same FWHM. As one can see, the simulated profile matches the observed one well at $\theta < 50'$, corresponding to about 0.7 times the virial radius, but is lower around the virial radius. Recently, Anbajagane et al. (2022) analyzed the SZ data of several hundred clusters and compared the results with simulations (Cui et al. 2018). They found that the average profile of the simulated clusters is also lower than the observation outside the virial radius. They suggested that the discrepancy might owe to the simulations not properly handling nonequilibrium effects, which might be significant because of the presence of accretion shocks and the low gas density in the outskirts of clusters. Another possibility is that our CSs do not model accurately the potentially important processes in the central part of the Coma cluster, such as the radio mode of AGN feedback. We will investigate these issues in future work.

Figure 10 shows the y -parameter map around the simulated Coma cluster. One can see a large, vertical edge located at $\alpha_J \sim 193.3$, which connects to a large edge extending from $(\alpha_J \sim 193.3, \delta_J \sim 27)$ to $(194.4, 26.6)$. This feature is outside the virial radius and can also be seen in the temperature and entropy maps shown in Figure 2. In the y -map of the Coma cluster obtained from PLANCK, one can also see a rapid decline of y at a similar location (see Figure 3 in Planck Collaboration et al. 2013c), and a diffuse radio relic is also observed there (see Figure 1 in Brown & Rudnick 2011). An

inspection of the simulation snapshots shown in Figure 2 reveals that the shock induced by the merger event at $z \sim 0.2$ produces this feature. In the y -profile of the simulated Coma shown in Figure 9, one can also see a significant drop at $r \sim 80'$, which roughly corresponds to the sharp edges seen in the y -map. Note that there is a substructure at a similar radius to the left of the main cluster, which might also contribute to the drop in the y -profile. The dashed curve in the left panel of Figure 9 shows the result excluding the region around the substructure. Clearly, the drop seen is not affected significantly by this substructure. All these demonstrate that the region around the Coma cluster provides an ideal avenue to detect and study the diffuse gas produced by accretion- and merger-driven shocks.

Recently, attempts have been made to detect the SZ effect produced by filaments. However, only a signal between two merging massive clusters was claimed to be detected (e.g., Planck Collaboration et al. 2013b; de Graaff et al. 2019). It is because the gas temperature and density in filaments is usually much lower than those in clusters, and a merging system can boost them. The SDSS Great Wall region contains many small and massive filamentary structures. Figure 11 shows the Compton y -map in the GW simulation. This map is obtained with the ParM method using a pixel size of $0'.5$. As shown in the Appendix, this method is reliable at $\log y > -8.5$ for such a pixel size. As one can see, the logarithm of the y -parameter in filaments is usually between -9 and -8 , much lower than the detection limit of current observations. The supercluster complex located at $\alpha_J \sim 184$ and $\delta_J \sim 4.5$ consists of three clusters with halo masses above $10^{14} h^{-1} M_\odot$ in a volume of $(10 h^{-1} \text{Mpc})^3$ and has a mean density about 17.2 times the cosmic mean density. A large fraction of this particular region is covered by pixels with $\log y > -8$, which may still be too

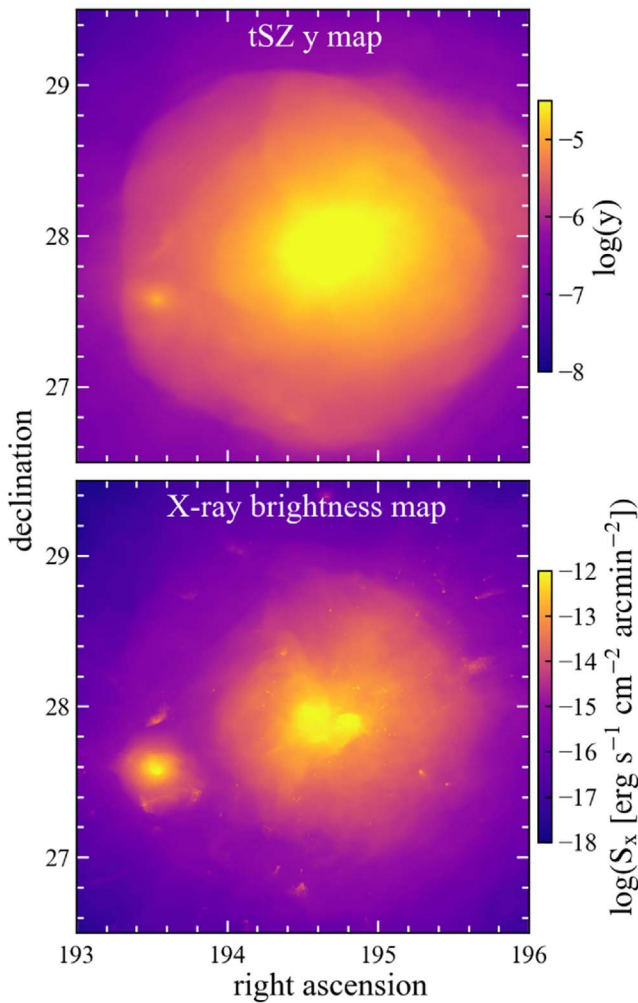


Figure 10. Maps of the y -parameter (upper panel) and X-ray (0.4–2 KeV band) brightness (lower panel) around the Coma cluster in the CM simulation.

weak to be detected directly by current observations. However, it may be possible to stack many pixels to have an average detection. Such a technique has already been applied successfully to detect the SZ effect produced by the CGM and the IGM (e.g., Lim et al. 2018a, 2018b).

Inspecting a series of snapshots of this region reveals a large number of violent merger events. These mergers are expected to produce strong shocks that could propagate outwards and heat the entire region. It may explain why this region has a larger y than other regions. In Figure 11, we also show y -maps around the three massive clusters in the supercluster complex. As one can see, there are many interfaces with sharp discontinuities around them. Some show sharp discontinuities at both small and large scales, reflecting the shocks induced by ongoing interactions in the recent merger histories of these systems. Our CSs will provide a unique data set to select interesting structures, such as filaments, for a stacking analysis, particularly when surveys of high sensitivity and resolution, such as CMB-S4 (Abazajian et al. 2016), become available.

Our CSs can also be used for cross-correlation investigations. For example, one can cross-correlate the simulated y map with observational data to study how well the model predictions match the observations. Owing to the weakness of the SZ effect, such cross-correlation analyses usually require large hydrodynamical simulations with a large sky coverage,

which are time consuming to run. However, as shown in Section 4, the local tidal strength λ_1 and the DM density ρ_m are tightly correlated with the gas temperature T and density ρ_{gas} , respectively. Thus, one can reconstruct the gas pressure field in the cosmic web. We can apply this method to the whole ELUCID N -body simulation in the SDSS region (Wang et al. 2016), or any other constrained cosmological N -body simulations, to predict the spatial distribution of electron pressure in the SDSS region. Assuming a simple relation between electron pressure and mass density, Lim et al. (2018b) obtained a model pressure map based on the ELUCID density field in the SDSS volume, and cross-correlated the model pressure with the Planck y -map to constrain the pressure–mass density relation over a large range of mass density. In particular they found that, at a given density, the y -parameter increases with the value of $|\lambda_1|$, as expected from its correlation with the gas temperature. The tight λ_1 – T relation found above suggests that such an analysis can be made much more powerful by incorporating the reconstructed local tidal strength. As shown in Figure 8, the typical scatter of the reconstructed pressure is about 0.25 dex over the density range relevant for filaments and sheets. This will provide a much improved criterion to select targets for stacking analyses.

5.2. X-Ray Emission

To calculate the X-ray surface brightness in a given band of $[E_1, E_2]$, we divide the sky of interest into many small pixels, each with a solid angle of $\Delta\Omega$. For each pixel, we calculate the surface brightness by summing up the contribution of all gas elements (cells in SphM method or particles in ParM method) in the pixel,

$$S_x(E_1, E_2) = \frac{1}{\Delta\Omega} \sum_i \frac{1}{4\pi D_L^2(z_i)} \times \int_{E_1(1+z_i)}^{E_2(1+z_i)} \Lambda(E, T_i, Z_i) n_{\text{H},i} n_{\text{e},i} dE, \quad (5)$$

where the emissivity, Λ , is taken from the look-up tables provided by the AtomDB (Foster & Heuer 2020, version 3.0.9)⁹; $n_{\text{H},i}$, $n_{\text{e},i}$, T_i , and Z_i are the number density of hydrogen ions, the number of free electrons, the temperature, and metallicity of the gas element i ; z_i and $D_L(z_i)$ are the cosmological redshift and luminosity distance of the gas element, respectively, calculated according to the distance of the element to Earth. The emissivity, Λ , provides both continuum and line emission under the assumption of collisional equilibrium, and all emission lines available in the chosen X-ray band are taken into account. Note that we exclude the contribution of gas elements within galaxies to the X-ray emission.

As in the last subsection, we use SphM to calculate the X-ray emission from the Coma cluster, and ParM for the GW simulation. A comparison between the two methods for X-rays is presented in the Appendix. Figure 9 shows the X-ray surface brightness profile of the simulated Coma cluster in the band of [0.7, 7] keV. For comparison, we also show the observational data from XMM-Newton (Mirakhor & Walker 2020, private communication). Similar to the y -parameter, our simulation is in good agreement with the observational data in the inner

⁹ <http://www.atomdb.org/index.php>

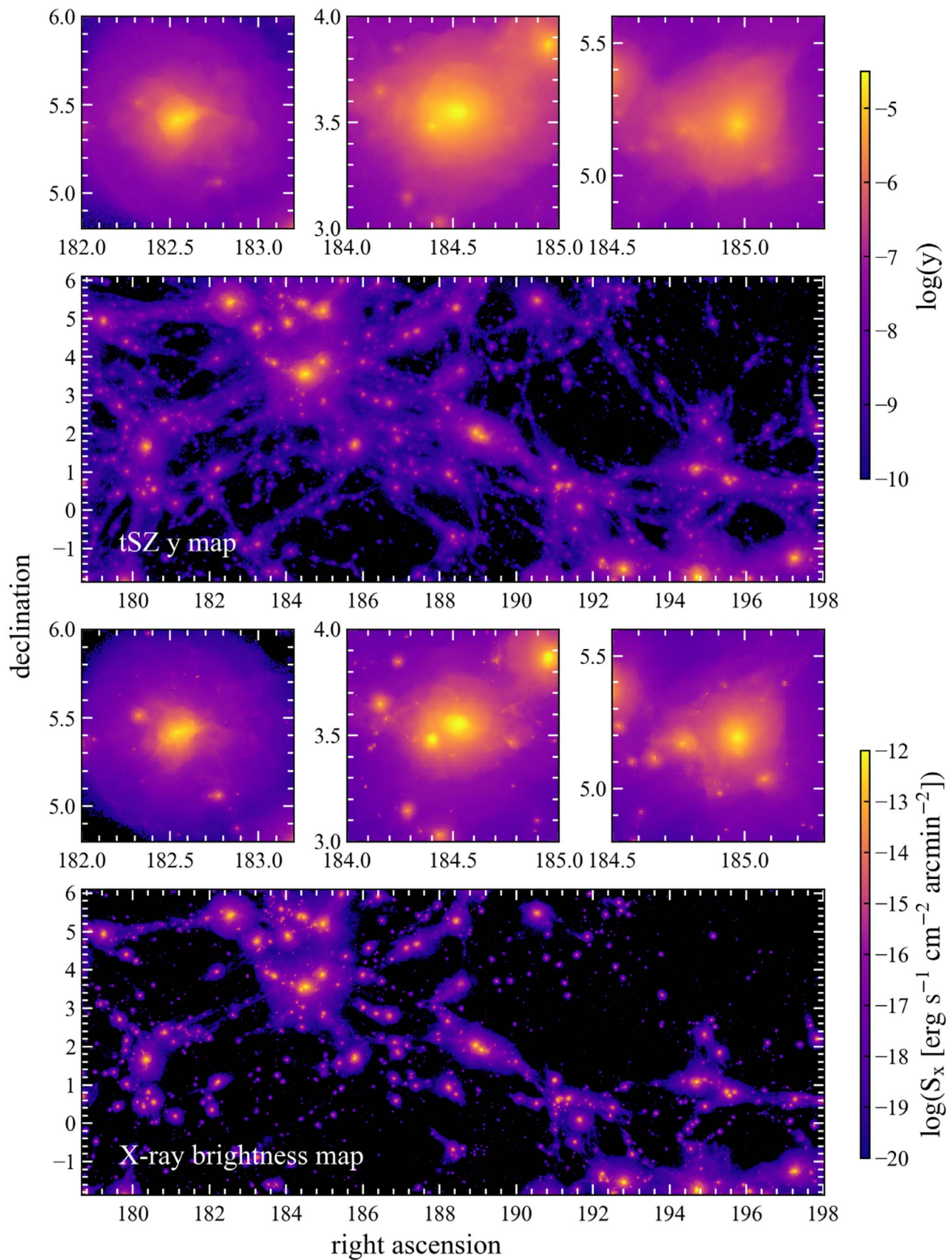


Figure 11. Maps of the y -parameter (the upper two rows) and X-ray (0.3–2 KeV band) brightness (the lower two rows) in the GW simulation. The first and third rows show the results for the three massive galaxy clusters in the supercluster complex located at $\alpha_j \sim 184$ and $\delta_j \sim 4.5$.

region from $10'$ to $40'$. In the most inner region, the simulation result shows a sharp rise. As shown in Huang et al. (2019), simulations using the PESPH technique may produce more dense gas in the central region of halos than the traditional SPH technique (e.g., Springel 2005). The small peaks on the X-ray profiles may also be caused by this artificial dense gas. In addition, our simulation does not include AGN feedback, and it is possible that AGN feedback can reduce the amount of high-density gas near the center. At $r > 40'$, the simulation prediction is systematically lower than the observation. The peak at $r \sim 70'$ in the simulation result is caused by the substructure discussed in the above subsection, and is significantly suppressed after the removal of the substructure (see the dashed curve).

Figure 10 shows the X-ray brightness map around the simulated Coma cluster. To see the inner structure more clearly and to compare with the eROSITA observations (Churazov et al. 2021), the brightness profile is obtained in the band of $[0.4, 2.0]$ KeV. One can see a bow-like sharp structure, more than 1° long, to the left of the cluster. There is another significant feature that is luminous and also has sharp edges at the center of the cluster. Interestingly, both features are seen in the eROSITA X-ray image of Churazov et al. (2021), who suggested that the feature to the left is likely produced by a shock while the one at the center is a contact discontinuity. Inspecting the formation history in our simulation indicates that the two features are actually generated by a merger that occurred more recently at $z \sim 0.2$.

In general, it is quite challenging to use X-ray emission to probe diffuse gas in the cosmic web, although attempts have been made to detect X-ray filaments associated with merging clusters (e.g., Sugawara et al. 2017; Reiprich et al. 2021). Figure 11 shows the predicted X-ray brightness map in the GW region, calculated in the band of $[0.3, 2.0]$ KeV. The pixel size of this map is $0'.5$. As shown in the Appendix, at such a pixel size, the ParM method is reliable at $\log S_x > -19$. The X-ray brightness drops very quickly as the distance to massive halos increases, because both the gas density and temperature decrease rapidly with the distance. The supercluster around $\alpha_j \sim 184$ and $\delta_j \sim 4.5$ might be a potential place to study diffuse gas in X-ray, as several relatively massive halos reside there. As a demonstration, we show the X-ray maps of the three most massive clusters in the region. As one can see, shock edges are also present around these clusters. The Cosmic Web Explorer, a proposed X-ray observatory, expected to reach $S_x = 10^{-18} \text{erg s}^{-1} \text{cm}^{-2} \text{arcmin}^{-2}$ in the band of $[0.3, 2.0]$ KeV (Simionescu et al. 2021), would make the X-ray emission from the diffuse gas predicted here easily detectable.

The good agreement between our simulation results and observational data in SZ and X-ray profiles, as well as in the locations of shock fronts and discontinuity features, in Coma cluster suggests that our simulation can reproduce the merging and heating histories of the Coma cluster reasonably well. Our simulations can thus provide a powerful avenue to interpret observational data. Our simulation results also have implications for AGN feedback. The fact that the predicted gas density and temperature profiles match the observed SZ and X-ray profiles at $10'$ to $40'$ suggests AGN feedback should not affect the gas significantly on these scales, because our current simulations do not include AGN feedback. Observations of both SZ effects and X-ray emissions of galaxy groups suggest that AGN feedback may be able to push a significant fraction of

baryonic gas out of their halos and heat the IGM in the cosmic web (e.g., Sun et al. 2009; Lovisari et al. 2015; Amodeo et al. 2021; Schaan et al. 2021). It is thus possible that the AGN feedback can dramatically change the X-ray and SZ predictions in regions of intermediate gas densities. The predicted X-ray and SZ signals in filaments shown in Figure 11 may thus be sensitive to the implementation of AGN feedback. Clearly, a detailed comparison between simulations of different AGN feedback and observational data are expected to provide important constraints on the underlying physical processes (see, e.g., Lim et al. 2018b). We will come back to this in a forthcoming paper.

5.3. Quasar Absorption Systems

Figure 12 shows HI and O VI column-density maps in the three sky regions. To obtain these column-density maps, we need the ionization states of gas particles. We use look-up tables from SPECEXBIN (Oppenheimer & Davé 2006; Ford et al. 2013) to find the ionization fraction for any gas particle in our simulations. These tables provide ionization fractions for various ions as functions of gas density, temperature, and redshift, obtained from CLOUDY (Ferland et al. 1998, version 08) assuming an optically thin slab of gas and collisional and photoionization equilibrium with a uniform Haardt & Madau (2012) background (hereafter HM12). For each sky region, we first select a large set of lines of sight that have a uniform distribution on the sky. We then divide each line of sight into many tiny intervals, each with a length equal to the gravitational softening length of $1.8 h^{-1}$ kpc. For each interval, we estimate the number density of the ion species in question from nearby gas particles using the SPH smoothing kernel.

Then we calculate the column density of an absorption line by integrating the density along the line of sight within the high-resolution region. As shown in Huang et al. (2019), assuming the HM12 UV background leads to a systematic overestimate of the HI column density. Following Davé et al. (2010); Huang et al. (2019) adjusted the HI optical depth by matching the evolution of Ly α flux decrements to observations and obtained a correction factor of 0.31. We thus multiply all HI column densities by the same factor. It should be pointed out, however, that the column density calculated in this way cannot be compared directly with the observed column density of individual absorption lines, as our analysis does not decompose the absorption into individual lines.

As one can see, high- N_{HI} gas, e.g., $\log N_{\text{HI}} > 15$, has a very compact distribution, while low- N_{HI} gas resides in filamentary structures. This is expected for low ionization energy ions, such as HI, which can be more easily destroyed by photoionization in gas of lower density. Almost all O VI absorption systems with $\log N_{\text{O VI}} > 13$ are produced by diffuse gas in filaments. These results indicate that low N_{HI} Ly α absorption and most of the O VI absorption with $\log N_{\text{O VI}} > 13$ are good tracers of the diffuse gas in filaments (see below; and Bradley et al. 2022). Compared to the O VI maps, the HI maps contain many smaller filamentary structures, indicating that low column-density Ly α absorption can be used to probe smaller filaments in the cosmic web, while O VI absorption mainly traces larger filaments. Halos residing in these thin filaments are usually very small, and thus have very low star formation efficiency (see, e.g., Zhang et al. 2021) and low metal abundances to produce significant O VI absorption. To verify this, we generated metallicity maps and found that the massive filaments are well

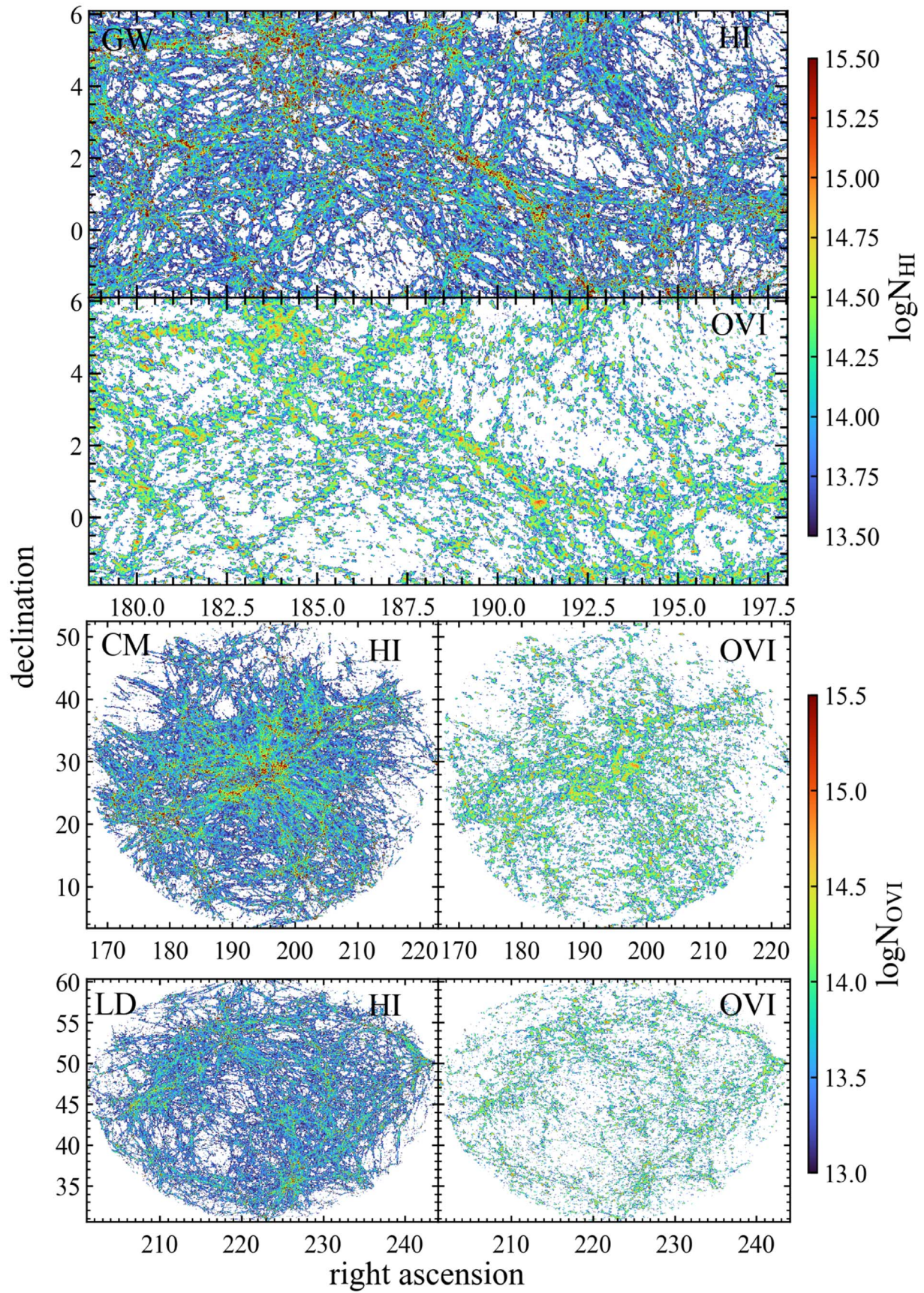


Figure 12. The HI and OVI column-density maps obtained from the three regions as labeled.

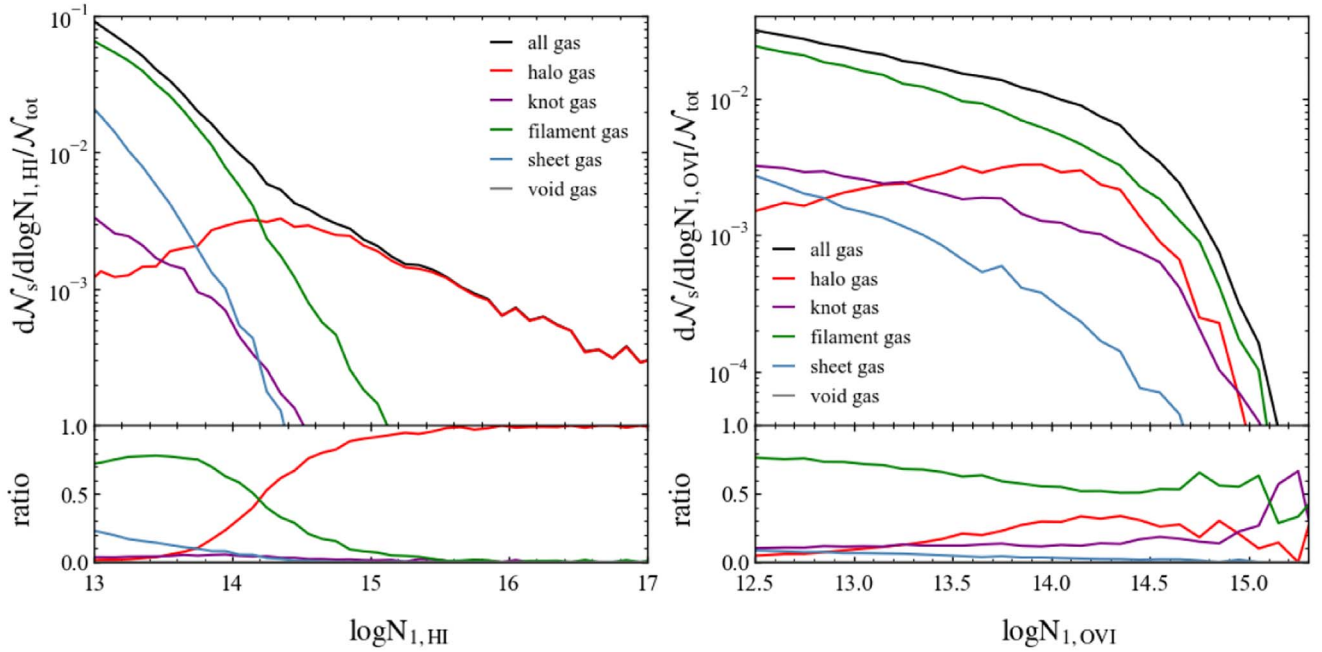


Figure 13. The upper panels show the distributions of H I (left) and O VI (right) column density obtained from the GW simulation. The column density is calculated by integrating the corresponding ion density over an interval of $1 h^{-1}\text{Mpc}$. Black lines show the results for all gas particles regardless of the types of the cosmic structure where they are located. The color-coded lines show the contributions of different components of the cosmic web. The contribution of the void gas is below the bottom of the plot. The lower panels show the ratio of the distributions of different components of cosmic web to all gas. See the text for details of the analysis.

represented, while the small filaments disappear, in the metallicity maps.

One can also see that the H I maps contain many more compact clumps of high column density than the O VI maps. As shown in Nelson et al. (2018), collisional ionization can effectively produce O VI at a temperature around $\log T \sim 5.5$ while photoionization is effective only in low-density regions. The gas in clumps with large N_{HI} usually has low temperature and high density, making it difficult to produce large amounts of O VI ions. Neither H I nor O VI absorption is a good tracer of massive halos, such as the Coma cluster, because the ICM is too hot to produce these ions. Interestingly, one can see many filaments in both H I and O VI around the Coma cluster, which are produced by relatively cold gas being accreted by the cluster. This is consistent with results from the Ly α absorber survey of the Coma cluster, which found that Ly α absorption tends to avoid the hot ICM (Yoon & Putman 2017).

More quantitative studies are required to understand which types of cosmic structure are traced well by these two types of absorption line systems. Figure 13 shows the distribution of the column density in different cosmic structures in the GW simulation. The column density in this figure is computed by integrating the density over an interval of $1 h^{-1}\text{Mpc}$ along a line of sight, instead of over the whole HIR. This column density is referred to as N_1 , and we sample the distribution of N_1 using a large number of random intervals. We classify each interval as halo, knot, filament, sheet, or void, according to the structure type that dominates the column density of the interval. The distribution of the column density produced by a given type of the cosmic web is defined as $dN_s/d \log N_i / N_{\text{tot}}$, where dN_s is the number of intervals with column density between $\log N_1$ and $\log N_1 + d \log N_1$ for the type in question, while N_{tot} is the total number of intervals regardless of structure types. We also show in the same figure the ratio of the distribution

function between the gas in a given cosmic web type and the total gas.

As one can see, the total $N_{1,\text{HI}}$ distribution is roughly a power law, similar to what is observed for quasar absorption line systems (see, e.g., Danforth et al. 2016). At $\log N_{1,\text{HI}} < 14.2$, the $N_{1,\text{HI}}$ distribution becomes steeper. The distribution functions for different types of the cosmic web show that, at $\log N_{1,\text{HI}} > 14.2$, the column density is dominated by gas in halos, while at $\log N_{1,\text{HI}} < 14.2$ it is dominated by diffuse gas associated with filaments. The contribution from the other three components is negligible in this column-density range. This is also seen in the column-density maps, where high column-density absorbers are compact and clumpy. Rudie et al. (2012, 2013) found that H I absorption systems with $\log N_{\text{HI}} > 15$ are spatially associated with galaxies, broadly consistent with our results that they are dominated by gas associated with halos.

The $N_{1,\text{OVI}}$ distribution is different from that for the H I. The total distribution is quite flat at $\log(N_{1,\text{OVI}}) < 14$ and declines rapidly at higher $N_{1,\text{OVI}}$, roughly consistent with observational results (see, e.g., Danforth et al. 2016). The decomposition into different cosmic web types clearly shows that the diffuse gas associated with filaments dominates over almost the entire column-density range plotted. More than 50% of the O VI absorption is associated with filaments, and the fraction increases to about 80% at the low $N_{1,\text{OVI}}$ end. Thus, O VI absorption is a promising probe of the diffuse gas component of the cosmic web. The Cosmic Origins Spectrograph (COS) Halos project (Tumlinson et al. 2011a; Werk et al. 2012; Tumlinson et al. 2013; Werk et al. 2013) detected many O VI absorption systems around galaxies in the local universe. This is not in conflict with our results. The COS-Halos project purposely selected quasar sightlines that are close to local galaxies, while our results shown above are obtained from sightlines that are uniformly distributed in the sky. In fact, our results show that halo and knot gas also have important contributions to high column-density systems with

$\log(N_{\text{I,O VI}}) > 13.5$. This is consistent with the COS-Halos project, which found that most of the detected O VI absorption systems have $\log N_{\text{O VI}} > 13.5$.

We also do the same analysis for the other two simulations. The shapes of the column-density distributions for both CM and LD are similar to those of the GW simulation. However, the amplitudes are smaller, as expected from the fact that the gas density in the GW region is higher than those in the other two (see Table 1). We also decompose the column-density distribution into different cosmic web components. At $\log N_{\text{I,HI}} > 14.2$, the distribution is dominated by halo gas, while at $\log N_{\text{I,HI}} < 14.2$, filament gas is the most important contributor. For O VI, on the other hand, diffuse gas in filaments dominates the column-density distribution over the column-density range considered here. These results are qualitatively the same as those for the GW simulation, and are not shown here. Note, however, that the cosmic web classification can be affected by the smoothing scale adopted in the calculation. To test the reliability of our results, we repeated our calculation using $R_s = 2 h^{-1} \text{Mpc}$, instead of $0.5 h^{-1} \text{Mpc}$, to make the classification. We also used different intervals to compute the column density. Our results are not affected significantly by these changes, and our basic conclusions remain the same.

We can also estimate the absorption column densities expected for real quasars in the three sky regions covered by our simulations. Based on the SDSS DR14 quasar catalog, we find 682, 13,711, and 6386 quasars with $0.2 < z_{\text{AGN}} < 0.9$ in regions covered by GW, CM, and LD, respectively. As shown in Danforth et al. (2016), the quasars in this redshift range can be used to investigate their absorption line systems. For each real quasar, we integrate the ion densities along its sightline to obtain the H I and O VI column densities. We find that there are 485, 8584, 4174 quasar sightlines with $\log N_{\text{HI}} > 13.5$ in the three regions respectively. Thus, more than 71%, 63%, and 65% of the quasars in the three regions are expected to have strong Ly α absorption lines. For a higher threshold of $\log N_{\text{HI}} > 14.2$, which selects absorption systems associated with halos, the numbers of quasar sightlines reduces to 128, 2150, and 784 in the three regions, respectively. For O VI, our simulations predict 301, 4855, 1801 quasar sightlines with $\log N_{\text{O VI}} > 13$ in their spectra, corresponding to 44%, 35%, and 28% of the quasars in the three regions, respectively. The fraction in the LD region is the lowest, while that in the GW the highest, as one expects given the difference in the mean mass density between the three regions. The numbers of quasar sightlines reduce to 96, 1536, and 522, respectively, when a higher threshold of $\log N_{\text{O VI}} > 14$ is used. It should be pointed out, however, other factors can affect the detection of absorption lines. For example, quasar spectra with a high signal-to-noise ratio are needed for the detection of weak absorption lines. Detailed investigations are required to select the proper targets to study quasar absorption line systems observationally.

As mentioned above, the column-density distributions shown above cannot be compared directly with the observation. In the future, we will model quasar absorption line systems in more detail, using SPECEXBIN (Oppenheimer & Davé 2006) to generate absorption profiles for various ions based on the predicted chemical abundance, ionization state, cosmic redshift, and peculiar velocity, and to identify individual absorption lines (see also Davé et al. 2010; Huang et al. 2019). This will allow us

to make direct comparisons with existing observational data in a statistical way. Furthermore, the results will also help us to identify the quasars that are the most promising targets for studying the properties of the gas associated with different components of the cosmic web. Given the large number of quasars available, most of the massive filament structures in the three regions can be sampled quite densely by quasar sightlines. The CSs can thus help to design observational strategies to maximize the observational efficiency. In addition, the full information about the accretion, heating, and enrichment histories provided by the simulations will open a new avenue to interpret the observational data of individual quasar sightlines in terms of underlying physical processes that determine the properties of the absorption gas.

6. Summary and Discussion

Studying the distribution and properties of gas in the cosmic web is a key step toward understanding galaxy evolution and formation, and in searching for *missing baryons* that are not observed in galaxies. In this paper, we use constrained hydrodynamic simulations to study the properties and evolution of the gas components in three representative regions of the local universe, the CM, the SDSS Great Wall, and a large low-density region at $z \sim 0.05$. The initial conditions of the three simulations are reconstructed from the ELUCID project. We use these simulations to demonstrate how the gas associated with different components of the cosmic web can be studied using the tSZ effect, X-ray emission, and quasar absorption line spectra. Our main results can be summarized as follows.

Our simulations reveal that cluster, filament, and low-density regions have very different evolutionary histories. The Coma cluster experienced several violent merger events during the past six billion years, while filament and low-density regions evolve more slowly. Gas in the cosmic web is heated by both adiabatic compression and shocks induced by mergers and accretion. About half of the WHIM resides in filaments, and the fraction does not change significantly among the three simulated regions.

We find a strong and tight correlation between the local tidal field strength (λ_1) and gas temperature (T). The relation consists of three power-law components: $\log T \propto \log \lambda_1$ at $\log \lambda_1 > 1$; $\log T \propto 2 \log \lambda_1$ at $-0.5 < \log \lambda_1 < 1$; and $\log T \propto \text{const.}$ at $\log \lambda_1 < -0.5$. This correlation, together with the tight correlation between the gas and total mass densities, can be used to predict the gas properties (density, temperature, and pressure) associated with the cosmic web predicted by a pure DM simulation, thereby providing a way to empirically model the cosmic gas and to interpret observational data.

Our CSs can reproduce the profiles of the Compton y -parameter and X-ray emission observed in the inner region of the Coma cluster. The simulations also produce, in maps of the y -parameter and X-ray, some discontinuity features that are seen in real data, suggesting that our simulations capture well the heating processes associated with the merging history of the Coma cluster. Thus, the CSs provide a new avenue to understand how the gas in and around the CM was affected by its formation history and by the baryonic physics assumed in the simulation.

We also investigate absorption line systems, such as Ly α and O VI, expected from our CSs. Ly α absorption systems with $\log N_{\text{HI}} > 14.2$ are mostly associated with gas within halos,

while most of the O VI absorption and low- N_{HI} absorption are produced by gas in filaments. O VI absorption lines are mainly connected to massive filaments, and small filaments, in particular those in low-density regions produce Ly α absorption systems with low column densities. Combined with the quasars observed behind the three simulated regions, our CSs can be used to select promising quasar candidates to explore the absorption of the gas associated with specific components of the cosmic web, e.g., massive filaments, thereby helping to understand how the spatial distribution and properties of the gas in the cosmic web are affected by formation processes and feedback processes.

The simulations presented here do not include AGN feedback. In the future, we will perform hydrodynamic simulations with different implementations of AGN feedback. As shown in previous studies using random phase simulations, AGN feedback can affect the gas component in the cosmic web in its thermodynamic and chemical properties, as well as its spatial distribution. Such feedback may thus significantly affect the predicted SZ, X-ray, and absorption line properties of the cosmic web. Detailed comparisons between CSs implementing different models of feedback and observational data in the same cosmic web thus provide an important avenue to discriminate feedback models.

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Appendix

In this paper, we adopted two methods to obtain SZ y -maps and X-ray brightness maps (Sections 5.1 and 5.2). The SphM method, in which gas properties are calculated on small cells with sizes less than the gravitational softening length using the SPH smoothing kernel, is the standard technique to obtain fluid quantities from SPH particles. However, this technique is very time consuming for a large simulation volume. The other method, ParM, which directly uses the properties of gas particles to estimate quantities on a given pixel, is much faster. However, the ParM method is valid only when the pixel size is larger than the SPH smoothing length. It is thus less accurate in lower-density regions.

In this Appendix, we compare the results obtained using the two methods. Here we select a small region in the GW simulation for the comparison. This region covers a large dynamical range in both y and S_x , which is ideal for testing the two methods. The maps of the two quantities obtained using the two methods, together with their fractional differences, are shown in Figure 14. Note that the maps for the ParM method use pixels of $0''.5$, the same as that shown in Figure 11, while the SphM maps have a much higher resolution. In calculating the difference maps, we use a pixel size of $0''.5$ for both SphM and ParM results.

As one can see, for both y and S_x , the two methods give almost identical results when their values are high. The differences become larger as y and S_x decrease. This is expected as the ParM method is not valid for low-density regions where the SPH smoothing kernel is larger than the pixel size and where both y and S_x are expected to be low. In Figure 15, we also show the difference between the two methods as a function of y_{SphM} and $S_{x,\text{SphM}}$, respectively. The fractional differences in y and S_x are less than 0.5 at $\log y_{\text{SphM}} > -8.5$ and $\log S_{x,\text{SphM}} > -19$, respectively, so that the results obtained from ParM are reliable for regions above these thresholds. A small number of pixels have $\log S_{x,\text{SphM}} > -19$ and yet a large difference between the two methods. As shown in Figure 14, these pixels are associated with the poor performance of ParM in handling small compact structures with sizes comparable to the pixel size. In general, the difference between ParM and SphM becomes smaller when the comparison uses a larger pixel size.

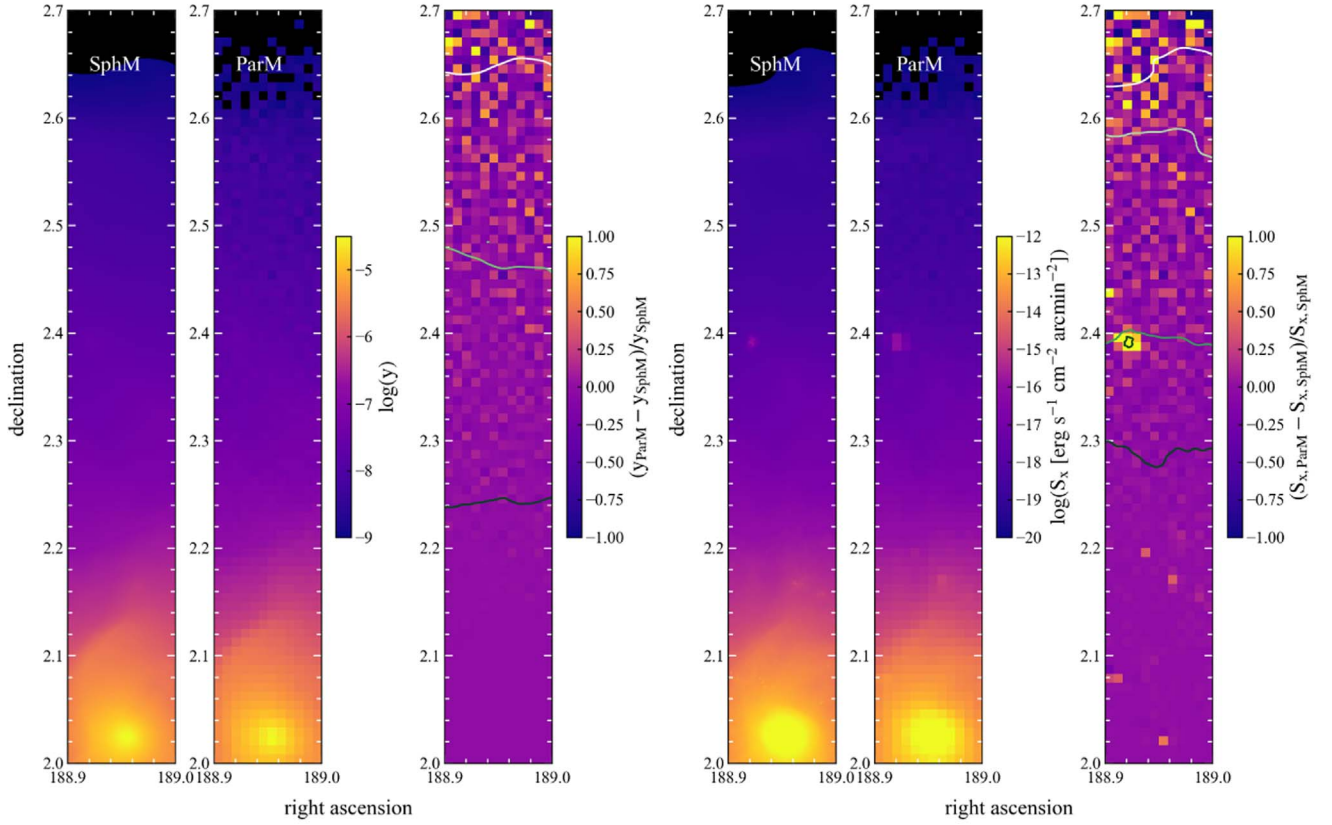


Figure 14. The left three panels show SZ y -maps of a small region in the GW simulation. The left one is obtained by using the SphM method, the middle one by using ParM, and the right one shows the difference between the two methods. The three contours in the difference map show $\log y_{\text{SphM}} = -7, -8, \text{ and } -9$, respectively. The right three panels show the results for the X-ray brightness, $S_{x,\text{SphM}}, S_{x,\text{ParM}}$, and their difference, respectively. The contours in the right panel show $\log S_{x,\text{SphM}} = -17, -18, -19, \text{ and } -20$, respectively.

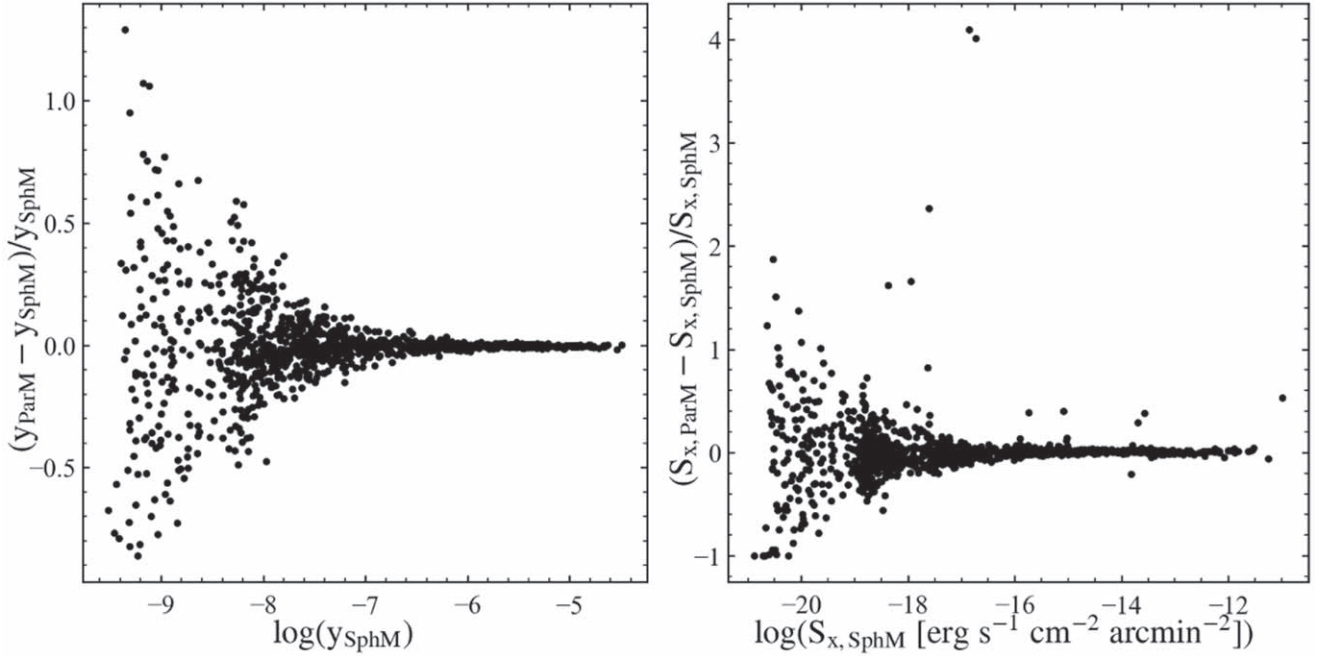


Figure 15. The left panel shows $(y_{\text{ParM}} - y_{\text{SphM}})/y_{\text{SphM}}$ as a function of $\log y_{\text{SphM}}$ for all pixels in the sky region shown in Figure 14. The right panel shows $(S_{x,\text{ParM}} - S_{x,\text{SphM}})/S_{x,\text{SphM}}$ as a function of $\log S_{x,\text{SphM}}$ for all pixels in the same sky region.

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