Planetesimal rings as the cause of the Solar System’s planetary architecture

Andre Izidoro, Rajdeep Dasgupta, Sean N. Raymond, Rogerio Deienno, Bertram Bitsch and Andrea Isella

Astronomical observations reveal that protoplanetary disks around young stars commonly have ring- and gap-like structures in their dust distributions. These features are associated with pressure bumps trapping dust particles at specific locations, which simulations show are ideal sites for planetesimal formation. Here we show that our Solar System may have formed from rings of planetesimals—created by pressure bumps—rather than a continuous disk. We model the gaseous disk phase assuming the existence of pressure bumps near the silicate sublimation line (at $T \sim 1,400$ K), water snowline (at $T \sim 170$ K) and CO snowline (at $T \sim 30$ K). Our simulations show that dust piles up at the bumps and forms up to three rings of planetesimals: a narrow ring near 1 au, a wide ring between ~3–4 au and ~10–20 au and a distant ring between ~20 au and ~45 au. We use a series of simulations to follow the evolution of the innermost ring and show how it can explain the orbital structure of the inner Solar System and provides a framework to explain the origins of isotopic signatures of Earth, Mars and different classes of meteorites. The central ring contains enough mass to explain the rapid growth of the giant planets’ cores. The outermost ring is consistent with dynamical models of Solar System evolution proposing that the early Solar System had a primordial planetesimal disk beyond the current orbit of Uranus.

The parent bodies of non-carbonaceous (NC) and carbonaceous (CC) chondrite meteorites are associated with classes of asteroids with overlapping orbital distributions and accretion ages, but different bulk chemical and isotopic compositions. NC and CC-like meteorites have water-poor and water-bearing compositions, which may suggest that they sample parent bodies that accreted inside and outside the snowline, respectively. They also show distinct stable-isotopic compositions for several volatile and non-volatile chemical elements. This isotopic ‘dichotomy’ implies that the millimetre-sized constituents of meteorites—pebbles—did not mix during the early accretion stages of planet formation, but their parent bodies did. The parent bodies of NC and CC meteorites must have accreted from distinct dust and pebble reservoirs that remained disconnected for 2–4 Myr (refs. 2,3). This is challenging given the predominant view that pebbles drift (mainly inward) quickly through the gas such that mixing of different reservoirs is unavoidable. Yet, after their assembly, asteroid-sized bodies must have dynamically mixed.

A possible missing ingredient comes from high-resolution ALMA observations revealing that rings and gaps are ubiquitous structures in protoplanetary disks. Ring-like structures suggest that pebbles have been concentrated in pressure bumps of the disk. At pressure bumps, the gas is locally accelerated, acting to trap drifting pebbles and increase the local solid–gas ratio, thus facilitating planetesimal formation. The existence of pressure bumps at different locations of protoplanetary disks is also supported by state-of-the-art disk models.

Here we propose a model of Solar System formation that accounts for the isotopic dichotomy and provides a link between the Sun’s natal disk and the structure of young disks around other stars. We first show that, if pressure bumps exist at important condensation fronts within the gaseous disk, planetesimals form in rings rather than a continuous disk. We use a suite of numerical simulations coupling different stages of planet formation—from dust coagulation to the final accretion phase of terrestrial planets—to demonstrate that the inner Solar System’s orbital architecture and meteorite data are consistent with the existence of up to three pressure bumps in the Sun’s natal disk. Our model accounts for the Solar System’s isotopic dichotomy, the formation of the terrestrial planets and the orbital distribution of NC and CC-like parent asteroids.

We model planetesimal formation by simulating dust coagulation, fragmentation and turbulent mixing in a young protoplanetary disk. We model dust evolution by solving the one-dimensional (1D) advection–diffusion equation for the dust column density. Micrometre-sized dust grains grow to millimetre- and centimetre-sized pebbles and start drifting inwards due to gas drag. As pebbles drift via drag they may also fragment. The local maximum pebble size depends on the local gas disk and grain properties, specifically the level of turbulence in the midplane, gas density and fragmentation velocity of icy and silicate pebbles. We solve the advection–diffusion equation for the largest (pebbles) and initial/smallest dust grains, assuming that the dust mass is mostly carried by largest grains. Pebbles inside the water snowline have silicate composition and smaller sizes relative to icy pebbles beyond the water snowline (Methods). This difference in sizes is imposed to reflect different threshold fragmentation velocities of ice-rich and silicate pebbles (see Supplementary Information for numerical tests of the importance of this parameter). In our model, silicate pebbles that reach regions of the gas disk hotter than 1,400 K are completely lost by sublimation. Ice-rich pebbles that cross the disk water snowline lose their ice, assumed to correspond to 50% of their bulk mass, releasing smaller silicate dust grains in the disk. Our gas disk model includes three (or two) pressure bumps. We model the gaseous protoplanetary disk by using radial power-law functions...
to represent the gas surface density ($\Sigma_{\text{gas}} \propto r^{-\beta}$, where we assume $x = 1$ or $x = 1.5$) and temperature profiles ($T_{\text{gas}} \propto r^{-\gamma}$), where we set $\beta \approx 0.7$–1.0. We mimic the presence of pressure bumps by rescaling the gas surface density by Gaussian and hyperbolic tangent functions at specific disk locations. The inner bump is associated with a transition in the gas disk’s viscosity due to thermal ionization of the gas at temperatures higher than $T_{\text{gas}} > 1,000$ K (refs. 14,15). It is located slightly outside the silicate sublimation line ($T_{\text{gas}} \sim 1,400$ K).

The central and outer pressure bumps are associated with the gas disk water snowline at $T_{\text{gas}} \sim 170$ K and CO snowline at $T_{\text{gas}} \sim 30$ K (refs. 14,17), respectively. The origin of pressure bumps at snowlines is usually associated with transition in the disk opacity due to different pebble sizes inside and outside the snowlines, and sublimation/condensation of grains at these locations11,12,18. Pressure bumps may also be associated with long-lived disk zonal flows25, although their location may not necessarily correlate with snowlines.

As expected,5,10 planetesimals naturally form at pressure bumps from pileups of drifting pebbles (Fig. 1a–c). In our model, planetesimal formation takes place if the local dust-to-gas midplane density ratio becomes higher than unity10, or if the pebble flux is high enough to promote pebble concentration via assistance of zonal flows in the gas disk4 (Extended Data Fig. 1). The pressure bump at the snowline works as an efficient barrier for icy pebbles drifting from beyond the snowline, disconnecting the inner and outer Solar System20 and promoting planetesimal formation. Only dust originally inside the snowline can contribute to planetesimal formation in the inner disk20. The pressure bump near the silicate sublimation line—at the interface between weakly and strongly ionized regions of the disk—is less efficient at trapping pebbles because they are smaller than icy pebbles and can more easily diffuse through and sublimate when they reach regions of the disk hotter than 1,400 K. Yet a fraction of pebbles is trapped at the inner bump and converted into planetesimals. As the disk evolves and cools, the pressure bumps move inward producing planetesimals from the outside in, forming three rings of planetesimals. The first planetesimals to form appear at $t < 100$ kyr, near 1.5 au, 8 au and 45 au. As the inner bump moves inwards, planetesimals are produced from 1.5 au to $\sim 0.7$ au (the inner ring). The bumps at the water and CO snowlines produce wider rings of planetesimals extending from $\sim 3$ au to 8 au (the central ring) and $\sim 20$ au to 45 au (the outer ring). Planetesimal formation in the inner disk ceases at $\leq 0.8$ Myr (see Supplementary Information for additional examples) because most pebbles originally in the inner parts of the disk are lost by inward radial drift. Planetesimal formation beyond the snowline proceeds while the pebble supply lasts and the gas is still around (see Supplementary Information for a discussion of the remaining mass in pebbles beyond the pressure bumps); we stop the simulation at 3 Myr. From $\sim 2$–3 Myr to the end of the gas phase (by $\sim 5$ Myr, growing and migrating giant planet cores would probably strongly influence planetesimal formation in the outer rings1, a process that we do not model here. The latest-formed planetesimals in the central ring form near $\sim 3$–5 au. Meteorite constraints forbid the inner and middle rings from overlapping2, assuming these to be associated with NC and CC planetesimals, respectively. We rule out scenarios where the water snowline moves inwards rapidly, reaching regions inside $\sim 1$–3 au early in the disk history (for example before Jupiter’s core formation; we discuss this issue in Supplementary Information).

Our simulations systematically produce two (Extended Data Fig. 2) or three rings of planetesimals (Fig. 1). The total mass and radial distribution of planetesimals in each ring depend on parameters such as the planetesimal formation efficiency ($\epsilon$), timing of pressure bump formation, gas viscosity in the magnetorotational active zone of the disk ($T_{\text{gas}} > 1,000$ K; refs. 14,15), level of turbulence in the gas disk midplane, initial dust-to-gas ratio ($Z_0$), gas disk width, and disk surface density and temperature profiles (Supplementary Information). In Fig. 1, the inner ring contains a total of $2.5 M_{\oplus}$ in planetesimals whereas the central ring carries $\sim 85 M_{\oplus}$. The outer ring has a total of $18 M_{\oplus}$ in planetesimals. If silicate pebbles are too small (for example when the level of turbulence in the disk midplane parameterized via $\alpha_t$ is sufficiently high, $\alpha_t \geq 10^{-4}$–$10^{-3}$) or the inner pressure bump is too weak (for example, when the region where $T_{\text{gas}} > 1,000$ K is not sufficiently ionized and the gas $\alpha$ viscosity at this location is $\alpha_{\text{gas}} \sim 10^{-3}$, where $\alpha_t$ corresponds to the viscosity when $T_{\text{gas}} < 1,000$ K; Methods), pebbles may not pile up sufficiently in the inner pressure bump and, consequently, the efficiency of planetesimal formation is markedly reduced in the inner ring. In contrast, if silicate pebbles are too large or if the inner bump is too strong, all pebbles inside the snowline may be converted into planetesimals at the inner pressure bump. The initial gas disk

![Fig. 1](VOL 6 | MARCH 2022 | 357-366 | www.nature.com/natureastronomy)
temperature also strongly controls the location of each planetesimal ring. In our preferred scenario, the initial disk temperature is typically ~1,000 K at ~1–1.3 au, which sets the initial location of the inner pressure bump. An initially colder disk where the gas temperature is 1,000 K at ~0.1 au would lead to a planetary system unlike the Solar System, if planetesimals can efficiently form in the inner pressure bump. The total mass in planetesimals in each ring also depends on the assumed planetesimal formation efficiency. Our preferred scenarios require planetesimal formation efficiencies varying from ~10^{-6} to ~10^{-3}, which is consistent with values assumed in previous studies.

We now model the evolution of the Solar System starting from a system with three rings of planetesimals similar to the one in Fig. 1. We consider constraints from the terrestrial and giant planets, the asteroid belt and meteorite measurements. We also use the inner Solar System's orbital architecture to further constrain the properties of the initial planetesimal rings. Our investigation proceeds from the inside out, from the growth of the terrestrial planets to the asteroid belt, giant planets and outer Solar System.

We model the growth of planetary embryos within the inner ring of planetesimals, assuming planetesimals to be originally born with a characteristic diameter of ~100 km (ref. 22). We compute planetary growth via pebble and planetesimal accretion using semi-analytical prescriptions calibrated from the results of N-body numerical simulations20,23. We feed our calculations with the evolving pebble flux and planetesimal surface density from our dust coagulation simulations (for example, Fig. 1) and follow the growth of planetesimals in simulations producing inner rings with different total mass in planetesimals. Our simulations show that in the inner ring where the total mass in planetesimals is ~2.5 \( M_\oplus \) (for example, Fig. 1) planetesimals grow to roughly Mars-mass planetary objects in ~1–3 Myr (Supplementary Information). These planetary embryos grow slowly and are unlikely to migrate substantially via type I migration26. In more massive disks the inner rings produce more planetesimals; for instance, a disk with a higher pebble flux (or higher planetesimal formation efficiency) produces an inner ring of 20 \( M_\oplus \) and forms Earth-mass planets in less than ~0.5 Myr (Supplementary Information). Such massive planets would migrate quickly24 and probably reach the innermost parts of the disk26,27. For this reason, simulations producing inner rings containing more than a few Earth masses in planetesimals are not consistent with the current Solar System. Our simulations also show that the growth of terrestrial planetary embryos in rings around ~1 au takes place via planetesimal accretion rather than pebble accretion, regardless of the ring’s total mass in planetesimals (Supplementary Information). The contribution from pebble accretion to the growth of protoplanetary embryos in the inner ring is negligible. This is because of the pressure bump at the snowline, which efficiently disconnects the outer and inner disk25. Pebbles in the inner disk (inside the snowline) are lost via radial drift before they can be efficiently accreted by growing planetesimals at ~1 au (ref. 20).

We performed N-body simulations of late-stage terrestrial accretion in the inner ring. The ring was assumed to extend from 0.7 au to 1.5 au and contain ~2.5 \( M_\oplus \) (as in Fig. 1) in planetesimals and planetary embryos following different surface density profiles, which accounts for potential different structures of the inner ring produced in our simulations. Figure 2 shows the terrestrial planets that formed after 200 Myr of integration produced in 80 simulations with slightly different initial conditions (Supplementary Information). The distribution of terrestrial planets in our simulations is in good agreement with the masses and orbits of the terrestrial planets in the Solar System. Our results are also consistent with previous formation models suggesting that the low mass of Mars relative to the Earth's mass reflects Mars's formation in a mass-depleted region20,27. Previous simulations25,26,29 had adopted ad hoc initial conditions, generally assuming that the asteroid belt was born mostly empty (the low-mass asteroid belt model). Here we have demonstrated the validity of this model from the earliest phases of the disk to planetesimal and terrestrial planet formation. Our model also addresses another Solar System mystery: the lack of planets inside Mercury's orbit. In our nominal simulation, the inner ring of planetesimals has a sharp edge at about 0.7 au due to sublimation of silicate pebbles at temperatures higher than 1,400 K preventing planetesimal formation inside ~0.7 au (refs. 20,21).

Our model sheds light on the isotopic dissimilarities between Earth and Mars. Martian meteorites show distinct isotopic signatures compared with Earth for several elements such as O, Ti, Ni, Mo, Nd, Cr and V22–24. Differences in mass-independent isotopic compositions between Earth and Mars suggest that these two planets accreted from broadly distinct materials from the Sun's natal disk. We investigate the source regions of protoplanetary objects constituting our final simulated planets in Fig. 2. Figure 3 shows that, in our best Solar System analogues (Supplementary Information), Earth and Venus analogues have broadly similar feeding zones, with most of their constituents (by mass) originating inside 1 au. In contrast, on average, only ~30% of the mass of our Mars analogues comes from inside 1 au, with some Mars analogues accreting entirely from material originally beyond 1.2 au. Although there is some variation in the feeding zones of Mars analogues, with some analogues also accreting mass from the inner parts of the inner ring26, these planets generally incorporate more mass from the outer parts of the inner ring (Extended Data Fig. 3 and Supplementary Information). We therefore propose that the origin of the distinct compositions of Earth and Mars is linked with how planetesimals form in the inner ring. Planetesimal formation in the inner ring starts at ~1.5 au. As the disk cools, and the inner bump moves towards the sun, the planetesimal production front also moves inwards. The moving inner bump is continuously fed by drifting pebbles from the outer parts of the inner disk while the supply lasts. For instance, at 0.04 Myr, planetesimals have started to form at ~1.5 au and the total mass in planetesimals between 2 and 5 au is ~12 \( M_\oplus \). These planetesimals will eventually reach the inner bump. Planetesimals forming at different locations
in the inner ring are made from pebbles arriving at different times at the bump, coming originally from different locations. Thus, our results suggest that the inner Solar System NC-like dust reservoir also had initially a subtle isotopic gradient that is reflected today in the bulk (isotopic) compositions of Earth and Mars40,41.

In our planetesimal formation simulations the main asteroid belt (from 1.8 to 3.2 au) is either born with very little mass in planetesimals of NC-like composition or completely empty (for example, in Fig. 1 the total mass in asteroidal planetesimals is \( \sim 10^{-4} M_\oplus \)—see Supplementary Information for parameter tests). While the present-day belt contains very little total mass, it has a complex orbital and compositional structure. The belt is dominated by two main groups, the C-type and S-type asteroids41.

A fraction of planetesimals from the terrestrial planet-forming region are scattered outward and implanted into the asteroid belt on stable orbits30,31. We associate these implanted planetesimals (as well as the small number of planetesimals that formed within the belt) with NC-like asteroids (the parent bodies of S-type/siliceous asteroids). We use our N-body simulations of the late stage of accretion of terrestrial planets to track the trapping efficiency of planetesimals from the inner ring into the belt over the first 200 Myr of the Solar System history. Planetesimals in the inner ring are scattered by massive planetary embryos on orbits that cross the asteroid belt region. To be captured into the belt, a scattered planetesimal must have its motion decoupled from the interacting embryo by an external perturbation that increases the pericentre of the object before it is ejected30,31. External perturbations may arise from other planetary embryos in the region or be due to orbital resonances with the giant planets in the asteroid belt.

Most implanted NC-like asteroids originated in the Mars region. In our N-body simulations of terrestrial accretion we imposed an initial planetesimal/embryo mass ratio of 30% that was constant throughout the inner disk. This choice is consistent with previous simulations30,31 and allows us to sample the trapping efficiency as a function of starting orbital radius anywhere within the inner ring. However, detailed calculations show that the planetesimal–embryo mass ratio should not be constant across the inner ring36. Rather, at the end of the gaseous disk phase the total mass in planetesimals near the inner edge of the terrestrial ring should be much smaller than that at the outer edge (compare black and coloured lines in the top panel of Fig. 4; see also Supplementary Information). This difference exists because planetary embryos in the inner parts of the ring grow faster due to shorter dynamical timescales, consequently consuming planetesimals faster than do embryos in the outer parts46. We account for this by rescaling our derived trapping efficiency (bottom panel of Fig. 4, black curve) by the radially dependent planetesimal–embryo mass ratio. While the non-normalized trapping efficiency is only a factor of 2 lower at 0.8 au compared with 1.4 au, the rescaled trapping efficiency is a factor of \(~10–100\) higher in the outer parts of the ring than in the inner parts (coloured lines in the bottom panel of Fig. 4).

The non-uniform implantation of planetesimals into the NC-like asteroid belt is consistent with Mars’s distinct chemical composition relative to that of Earth. S-type asteroids are linked with ordinary chondrite meteorites.44 Martian meteorites and ordinary chondrites are similar in isotopic signatures of Cr, Ti, O and V, and Mars is thought to be mostly made of material akin to ordinary chondrites35,46. This compositional similarity is naturally explained by the high implantation efficiency of planetesimals from the Martian region relative to the Earth region (for example \(~1.2–1\) au), and the different feeding zones of the two planets (Fig. 3). Enstatite chondrites, associated with the less abundant E-type asteroids3, may sample planetesimals originally inside \(~1\) au. This is consistent with isotopic models for Earth’s formation suggesting enstatite chondrites as the major constituent37,38.

The C-type asteroids were probably implanted from the giant planet-forming region during the gaseous disk phase. If we assume that pebbles and planetesimals beyond the snowline have CC-like compositions, we can compare their orbital evolution with that of present-day carbonaceous asteroids (C-type), which are broadly distributed across the entire main asteroid belt between 1.8 au and 3.2 au (ref. 1). CC-like planetesimals formed exterior to 3–4 au (Fig. 1), and must have interacted with the gas giant planets36,37. As Jupiter and Saturn’s cores grew and migrated, they scattered nearby planetesimals onto eccentric orbits; a fraction were implanted into the asteroid belt due to the effects of gas drag in the disk36. The implantation efficiency of planetesimals from the central region into the belt is a steep function of orbital distance, with the innermost planetesimals in the central ring having the highest probability of being implanted (Fig. 1d). This implies that late-formed planetesimals, typically born \(~2–3\) Myr after the start of our simulations, have the highest implantation probabilities. In our model, these correspond to the parent bodies of CC-like chondrites whose accretion ages are estimated to be at least 2 Myr after the formation of the calcium–aluminium inclusions5. Planetesimals beyond the snowline formed during the first 1 Myr are also implanted into the asteroid belt but with much smaller probabilities. These objects are broadly consistent with the formation ages of parent bodies of iron meteorites5. Note that the implantation of planetesimals from the middle ring into the belt and the trapping of planetesimals from the inner ring into the belt region are dictated by different mechanisms. Whereas the implantation of planetesimals from the central ring is assisted by gas drag, the trapping of planetesimals from the inner ring is virtually independent of gas drag and is due to gravitational interactions.

Our model can match the inner Solar System’s chemical and isotopic constraints. In our simulations, planetesimals in the inner ring form during the first \(~0.5–1\) Myr of the Sun’s disk lifetime (Supplementary Information). The decay of short-lived radioactive isotopes such as \(^{26}\)Al is expected to trigger melting and differentiation.
Articles

Fig. 4 | Planetesimal implantation efficiency into the asteroid belt for different inner ring configurations. Top: planetesimal-embryo disk mass ratio as a function of orbital distance. Blue, yellow, green and purple show the planetesimal-embryo disk mass ratio at 5 Myr (timing of the gas disk dispersal), in simulations modelling the growth of planetesimal in rings around 1 au. In all cases, the ring extends from ~0.7 to ~1.5 au, as in Fig. 1, and starts with ~2.5 $M_\oplus$ in 100-km-sized (diameter) planetesimals. Planetary objects are defined as objects more massive than the Moon. Colour-coded lines also show rings with different initial planetesimal surface density profiles, proportional to $r^{-1}$ (green), $r^{-3}$ (blue), $r^{-2}$ (purple), $r^{-1.5}$ (pink) and ($-200(r(\text{au})-1)^2+24)$ g cm$^{-2}$ (yellow). The black curve shows the disk mass ratio in our reference simulation. Bottom: calibrated planetesimal trapping efficiency into the asteroid belt as a function of orbital distance. Colours represent different rings as in the top panel. The black line-symbols show the reference planetesimal trapping efficiency into the asteroid belt derived from our simulations modelling the late stage of accretion of terrestrial planets, after 200 Myr of integration.

drules in the disk, if it interacts with a population of planetesimals existing in the belt (for example, Fig. 1; or beyond its orbit in the case of CC planetesimals). Chondrules generated via different processes may either reaccrete onto first-generation planetesimals or trigger the formation of a second generation of planetesimals if the local solid-to-gas ratio becomes sufficiently large to promote planetesimal formation at the very end of the gas disk lifetime (see Supplementary Information for estimates of dust production in our simulations modelling the growth of planetesimals to planetary embryos). Our results suggest that ordinary chondrites in reality most probably sample differentiated, partially differentiated and perhaps even undifferentiated parent bodies.

A model with multiple rings of planetesimals also matches the outer Solar System. The integrated mass in the giant planets’ cores is ~60–100 $M_\oplus$ (ref. [7]). The ring of planetesimals associated with the water snowline, extending from 4 to ~10 au (Fig. 1), typically contains ~40–100 $M_\oplus$ (Supplementary Information), and ~0–100 $M_\oplus$ in leftover pebbles. Massive cores form quickly by pebble accretion onto the most massive planetesimals. Simulations that match the ice giants’ mass distribution invoke a phase of giant collisions between ~5 $M_\oplus$ cores whose inward migration was blocked by the near fully grown Jupiter and Saturn. We can envision the growth of the ice giant cores either from the outer parts of the middle ring of planetesimals or the inner parts of the outer ring of planetesimals.

It is well accepted that the giant planets underwent a dynamical instability (see Supplementary Information for a discussion of our model in the context of giant planet migration and Solar System evolution), which spread out and excited the giant planets’ orbits, and can explain the orbital distributions of many small-body populations. The instability may have been triggered by dynamical interactions between the giant planets, whose orbital configuration was compact at early times, and an outer disk of planetesimals. MATCHING THE GIANT PLANETS’ ORBITS REQUIRES A PRIMORDIAL OUTER DISK CONTAINING ~10–30 $M_\oplus$ (ref. [11]). The outer planetesimal disk in Fig. 1 contains ~18 $M_\oplus$ and fits nicely with dynamical models.

The cold classical Kuiper belt is a population of objects on low-eccentricity, low-inclination orbits extending from ~42 au out to ~45 au. The cold classicaKBOs are probably never strongly scattered by the planets and may thus represent the outermost planetesimals that formed around the Sun. While all Kuiper-belt objects (KBOs) show red colours, the cold classicaKBOs show far redder colours than other KBOs with similar orbital radii but with larger eccentricities and inclinations. It has been proposed that KBO colours correlate with their formation distance from the Sun. Our multiple-ring model naturally explains the colour dichotomy, as the cold classicaKBOs would represent remnants from the distant parts of the outermost planetesimal ring whereas more dynamically excited KBOs originated in the outer parts of the central ring or the inner parts of the outermost ring.

It is legitimate to wonder how generically applicable our model is to the formation of other planetary systems. Observations and statistical analysis suggest that planets with sizes between those of Earth and Neptune (1 and 4 $R_\oplus$) are common around other stars. These planets are typically referred to as super-Earths. Super-Earths with orbital periods shorter than 100 d have been measured to orbit at least 30% of the Sun-like stars. Mercury’s orbital period is ~88 d, yet no planet inside Mercury’s orbit exists. Why? We have demonstrated that our simulations can naturally lead to the formation of two classes of planetary systems. In systems where a strong pressure bump forms early at the snowline, pebbles from the outer disk are trapped at the bump and prevented from drifting inwards to the inner disk. This efficient disconnection of the inner and outer system can explain the Solar System isotopic dichotomy as well as the lack of massive planets (for example super-Earths) in the terrestrial region and potentially inside Mercury’s orbit. In this particular scenario, planetary embryos growing in the terrestrial region
grow at most to (a few) Mars-mass planetary embryos and avoid large-scale radial migration, forming a system of terrestrial planets similar to those in the solar system. On the other hand, in systems where the pressure bump forms late (Extended Data Fig. 4) or is not as strong (more leaky), the inner system is invaded by tens of hundreds of Earth masses in pebbles from the outer disk. In such disks, Earth-mass (or more massive) planetary embryos are likely to form rapidly near 1 au and undergo large-scale radial migration, eventually reaching the inner edge of the disk and forming systems of hot super-Earths. The fundamental question that emerges is why some systems form efficient pressure bumps (for example, the Solar System) whereas others do not (for example, super-Earth systems).

We propose that this may be linked to intrinsic characteristics of the protoplanetary disk, such as the level of turbulence (viscosity) in the disk controlling pebble sizes ($a_t$), or the existence of favourable conditions for the early formation of giant planets (for example Jupiter) at specific locations of the disk, as at the disk water snowline. Pressure bumps in gaseous disks with smaller pebbles may be far more leakier than those in disks with larger ones, potentially allowing a lot of mass in pebbles from the outer disk to be delivered to the terrestrial region. In our Solar System, Jupiter’s early core formation may have also played a decisive role in preventing this from happening. If a giant planet core promptly forms in the pressure bump at the snowline, it may induce the formation of another pressure bump beyond its orbit. The Solar System architecture, which seems to be an unusual outcome of planet formation, probably reflects all these conditions.

Our model demonstrates that the present-day Solar System could have formed from three rings of planetesimals (Fig. 5). This diverges from standard models that assume a continuous disk of planetesimals, and is reminiscent of the ring-like structures observed in disks around young stars. Our model can simultaneously explain the orbits and masses of terrestrial planets and the distribution of different types of asteroid in the main asteroid belt, and link different classes of meteorites to the main building blocks of Earth and Mars. The terrestrial planets formed from a narrow ring in the terrestrial region, which naturally accounts for the lack of planets inside Mercury’s orbit and the low mass of Mars. The gas giant planets formed from a wide central ring located beyond the asteroid belt, probably via planetesimals and pebble accretion. The asteroid belt was never much more massive than it is today. The outermost planetesimal ring located beyond the current orbit of Uranus was sculpted during the giant planets’ growth and their subsequent dynamical evolution, producing the current Kuiper belt.

**Methods**

Our model couples different stages of planet formation, from dust evolution in a young gaseous disk to the final stage of accretion of terrestrial planets. We have performed simulations modelling the following.

- Dust evolution (growth and fragmentation) in a gaseous disk
- Planetesimal formation
- Growth from planetesimals to planetary embryos in rings around 1 au
- The late stage of accretion of terrestrial planets—namely, growth from planetary embryos to planets
- Asteroid implantation from inside out during the terrestrial planet formation
- We next describe the different ingredients of our model individually.

**Gas disk model.** We model the radial distribution of gas and disk temperature using power-law profiles. Gas disk dissipation and cooling are mimicked via exponential decay timescales. Although simplistic, our approach allows us to easily disentangle the effects of different parameters of the model. In nominal simulations, our 1D underlying protoplanetary disk extends from 0.1 au to ~120 au and is represented by a simple power-law disk with radial profile given as

\[
\Sigma_{\text{gas}}(r) = 1.70 \left( \frac{r}{1 \text{ au}} \right)^{-1} \exp \left( -\frac{r}{r_{\text{ref}}} \right) \frac{\text{g cm}^{-2}}{\text{cm}^{-2}}.
\]

In equation (1), $r$ is the heliocentric distance. Following previous studies, we mimic the presence of the pressure bumps in the disk by rescaling the original gas disk profile given by equation (1) with Gaussian and hyperbolic tangent functions. Each of these functions accounts for one of the pressure bumps in the disk. Our gas disk radial profile reads

\[
\Sigma_{\text{gas}}(r) = \Sigma_{\text{gas}}(r) \exp \left[ -A_{\text{CO}} \exp \left( -\left( \frac{r - r_{\text{cold}}}{r_{\text{cold}}/10^5} \right)^2 \right) \right] \exp \left[ -A_{\text{CO}} \exp \left( -\left( \frac{r - r_{\text{cold}}}{r_{\text{cold}}/10^5} \right)^2 \right) \right] \alpha_r.
\]

where

\[
S(r) = 0.5 \left( \alpha_{\text{gas}} - \alpha_r \right) \left( 1 - \tanh \left( \frac{T_{\text{fin}} - T_{\text{gas}}}{50 \text{ K}} \right) \right) + \alpha_r.
\]

In equation (2), $\alpha_{\text{gas}}$, $r_{\text{cold}}$, $A_{\text{CO}}$, $w_{\text{cold}}$, $w_{\text{CO}}$, and $H_{\text{CO}}$ are free parameters used to represent the pressure bump locations, amplitudes and widths. $H_{\text{CO}}$ and $H_{\text{CO}}$ are the gas disk scale heights at the water ($T_{\text{water}} = 170$ K) and CO ($T_{\text{CO}} = 30$ K) snowlines, respectively. Following previous studies, we set $A_{\text{CO}} = 0.5$ and $w_{\text{CO}}(r_{\text{cold}}) = H_{\text{CO}}(r_{\text{cold}})$. We choose for the pressure bump at the snowline a configuration that provides an efficient disconnection of the inner and outer Solar System planet reservoirs, as constrained by the Solar System isotopic dichotomy. For simplicity, we assume an equivalent configuration for the pressure bump at the CO snowline, that is, $A_{\text{CO}} = 0.5$ and $w_{\text{CO}}(r_{\text{cold}}) = H_{\text{CO}}(r_{\text{cold}})$. $\alpha_r$ is the gas disk viscosity in regions of the disk where $T_{\text{gas}} < T_{\text{water}} = 1,000$ K. In
our model, \( T_{\text{gas}} \) defines the threshold temperature for thermal ionization of the gas disk\(^6\). We set \( \alpha = 10^{-5} \) in all our simulations\(^7\). We assume that in regions of the disk where \( T_{\text{gas}} > T_{\text{sam}} = 1,000 \text{~K} \) the disk viscosity increases due to the presence of magnetorotational instabilities fostered by thermal ionization of the gas\(^8,9\). Previous studies have assumed that the gas disk viscosity in strongly ionized regions of the disk may be a factor of \( 10-100 \) larger than in weakly ionized regions\(^10\). In our nominal simulations, we assume a conservative value corresponding to \( \alpha_{\text{sam}} = 5 \alpha_0 \). At the transition between the high- and low-viscosity regions of the disk, the gas surface density drops to keep the gas accretion rate constant as a function of radius. This creates a pressure bump in the disk slightly outside the silicate sublimation line of the disk\(^11\). Note that \( \alpha_{\text{sam}} \) and \( \alpha_0 \) set the shape and strength of the inner pressure bump. We explore the effects of this parameter in our model in simulations presented in Supplementary Information.

The initial gas disk temperature is modelled as a simple power-law disk as

\[
T_{\text{gas}}(r) = 1,200 \left( \frac{r}{1 \text{~au}} \right)^{-2} \text{~K},
\]

where the power-law index \( p \) is assumed to be equal to either 0.7 or 1 (ref. \(^7\)). That is, we test our model against two disk temperature profiles. For \( p = 0.7 \), the temperature profile places the CO snowline initially at \(-200 \text{~au}\). As our nominal disk extends up to \(-120 \text{~au}\), simulations with \( p = 0.7 \) contain only two pressure bumps. In cases where \( p = 1.0 \), the pressure bump at the CO snowline is initially at \(-41 \text{~au}\).

Our initial disk temperature profiles are broadly consistent with models of disk formation from the collapse of the progenitor molecular core\(^12\). The time zero of our simulations probably corresponds to the first \( \sim 10-100 \text{~kyr} \) from the beginning of the collapse, when the disk has already expanded and starts to cool off in the viscous energy\(^7,13\). In all simulations, we assume that the gas disk dissipates following an exponential decay with \( e \)-fold timescale of \( 2 \text{~Myr} \). The gas disk temperature also decays exponentially with time—but with different timescales as the disk evolves—to better mimic the evolution of the disk temperature seen in simulations modelling disk evolution\(^15\). In Fig. 1, where \( p = 1.0 \), the disk temperature starts to drop with a short \( e \)-fold timescale of \( 0.75 \text{~Myr} \) at \( 0.05 \text{~Myr} \). From \( 0.5 \text{~Myr} \) to \( 2 \text{~Myr} \), we increase this timescale to \( 5 \text{~Myr} \) to roughly mimic the evolution of the snowline in more sophisticated disk models\(^16\). From \( 2 \text{~Myr} \) to \( 3 \text{~Myr} \), we assume that the disk temperature is not evolving with time. This approach is used to mimic Jupiter’s core formation in the disk (see also Supplementary Information for parameter tests). Once Jupiter’s core forms, perhaps near/at the snowline pressure bump\(^16\), it prevents pebbles from beyond its orbits drifting inwards, and may even heat up the disk, shifting the location of the snowline further out\(^16\). We do not include these effects in this work, but they are not expected to impact qualitatively our main conclusions.

The final position of the snowline in our disks is broadly consistent with the results of more sophisticated disk models\(^16\). Note that our disk is initially much hotter than that considered in our previous study modelling the effects of a pressure bump on the formation of the Solar System terrestrial planets\(^17\). In ref. \(^7\), the disk snowline was initially at \( 5 \text{~au} \) and the transition in the disk viscosity due to thermal ionization would be at \(-0.1 \text{~au} \). The disk models invoked here and that of Izidoro et al.\(^18\) represent two possible end-member scenarios. The main advantage of our new model is that it provides a simple solution to one of major problems of the scenario of Izidoro et al.\(^18\): our new model naturally accounts for the lack of planets inside Mercury’s orbit in the Solar System, a problem that remained unsolved in ref. \(^7\).

**Dust evolution calculations.** We model coagulation, fragmentation, drift and turbulent mixing of dust grains in a 1D gaseous disk. Our dust evolution code is tests\(^19\). Once Jupiter’s core forms, perhaps near/at the snowline pressure bump\(^6\), it prevents pebbles from beyond its orbits drifting inwards, and may even heat up the disk, shifting the location of the snowline further out\(^16\). We do not include these effects in this work, but they are not expected to impact qualitatively our main conclusions.

The final position of the snowline in our disks is broadly consistent with the results of more sophisticated disk models\(^16\). Note that our disk is initially much hotter than that considered in our previous study modelling the effects of a pressure bump on the formation of the Solar System terrestrial planets\(^17\). In ref. \(^7\), the disk snowline was initially at \( 5 \text{~au} \) and the transition in the disk viscosity due to thermal ionization would be at \(-0.1 \text{~au} \). The disk models invoked here and that of Izidoro et al.\(^18\) represent two possible end-member scenarios. The main advantage of our new model is that it provides a simple solution to one of major problems of the scenario of Izidoro et al.\(^18\): our new model naturally accounts for the lack of planets inside Mercury’s orbit in the Solar System, a problem that remained unsolved in ref. \(^7\).

In our simulations, we explored different values for \( Z_0 \) ranging from 0.25% to 1.5% (ref. \(^8\)). To model dust evolution, we solve the 1D advection–diffusion equation for the column dust density\(^20\)

\[
\frac{\partial \Sigma_{\text{dust}}}{\partial t} + \frac{1}{r} \frac{\partial}{\partial r} \left( \sqrt{\Sigma_{\text{dust}} D_{\text{dust}}} \frac{\partial \Sigma_{\text{dust}}}{\partial r} \right) = \frac{-\frac{1}{2} \Sigma_{\text{gas}}}{\partial r} \frac{\partial \Sigma_{\text{gas}}}{\partial r} \frac{\Sigma_{\text{gas}}}{\partial r} = \frac{\Sigma_{\text{gas}}}{\partial r} \frac{\partial \Sigma_{\text{gas}}}{\partial r} \frac{\Sigma_{\text{gas}}}{\partial r}
\]

where \( \Sigma_{\text{gas}} \) represents the local gas-to-dust ratio, which evolves with time. The pressure support parameter is calculated as

\[
\eta = -\frac{1}{2} \frac{c_s}{v_k} \left( \frac{\partial \ln P}{\partial \ln T} \right) \frac{\Delta}{P}
\]

where \( v_k \) is the Keplerian velocity. The nominal gas radial velocity is calculated as\(^21\)

\[
v_{r,\text{gas}} \approx -\frac{c_s}{r}
\]

where the gas disk viscosity\(^22\) is \( \nu = \alpha H_{\text{gas}} \) (ref. \(^7\)). \( c_s \) represents the sound speed, \( P \) is the midplane gas disk pressure calculated in the isothermal limit and \( H_{\text{gas}} \) is the disk scale height, derived from the temperature profile assuming that the gas disk is in vertical hydrostatic equilibrium. \( H_{\text{gas}} = H_{\text{disk}} \) is the gas disk aspect ratio. We include the effects of back-reaction of the dust on the gas disk as\(^23\)

\[
v_{r,\text{gas}} = \frac{StZ^2}{\partial r} \left[ 1 + \frac{(1 + Z^2)^2}{\partial r} + \frac{(1 + Z^2)^2}{\partial r} \right] v_{r,\text{gas}} \]

The critical pebble flux given in equation (8) is defined as\(^24\)

\[
\nu_{\text{crit}} \sim \frac{\Sigma_{\text{dust}}}{\partial r} \frac{\partial \Sigma_{\text{gas}}}{\partial r} \frac{\Sigma_{\text{gas}}}{\partial r} \frac{\Sigma_{\text{gas}}}{\partial r} \frac{\Sigma_{\text{gas}}}{\partial r}
\]

Water ice particles that drift inwards and cross the water snowline are assumed to sublimate, losing 50% of their mass, corresponding to the ice component. This effect is particularly important when the pressure bump at the water snowline is not present from the beginning of the simulation, and a substantial mass in pebbles from beyond the water ice line drifts inside it before the bump forms. Silicate pebbles in the inner disk are assumed to sublimate when crossing the silicate sublimation line (\( T_{\text{sam}} = 1,000 \text{~K} \)). In our simulations, we use a radial log grid with 400 cells, which ensure numerical convergence\(^25\). Following previous studies\(^26\), our model assumes that the threshold fragmentation velocity of ice pebbles beyond the snowline is \( 10 \text{~m} \text{s}^{-1} \), whereas silicate pebbles fragment at a velocity of \( 1 \text{~m} \text{s}^{-1} \). We use a smoothing function to model the transition in the threshold velocity at the snowline\(^26\). The turbulence level at the disk midplane is represented by \( \alpha \) and the gas disk viscosity by \( \alpha_0 \) as adopted in similar studies\(^27\). In our simulations, we assume that \( \alpha = \alpha_0 \Delta \), where \( \Delta \) is a simple dimensionless free parameter, which we assumed to range from 1 to 40 (ref. \(^8\)). We discuss the impact of these parameters on our model in Supplementary Information. Note that, for these levels of turbulence in the disk midplane, planetesimals are unlikely to form during the disk build-up\(^28\). Therefore, our simulations are envisioned to correspond to the timing when the gas in-fall from the molecular cloud has ceased, and the disk is fully formed. In this work, we do not account for the recondensation of vapour diffusing back to regions beyond the sublimation line when drifting pebbles cross a giving sublimation front\(^29\) because we do not solve the advection–diffusion equation of the gas. We also solve the advection–diffusion equation for the dust component considering two dust-size populations rather than a ‘continuum’ of grain sizes. We do not expect these assumptions to impact the broad qualitative results of our paper.
The collapsing dust mass \( m_c \) is set to
\[ m_c = \frac{4}{3} \rho_H \pi l_H. \tag{14} \]
\( r_c \) represents the characteristic lifetime of traps due to zonal flows that promote pebble concentration until they can collapse. We set \( r_c = 100 \) Myr (refs. [1–3], this letter) which represents the Hill density defined as
\[ \rho_H = \frac{9 M_\oplus}{\pi r_H^3}. \tag{15} \]
\( M_\oplus \) is the central star mass. The critical length scale can be derived by equating the diffusion and collapse timescales of the clump, which yields
\[ l_H = \frac{2}{3} \sqrt{\frac{\alpha}{\omega}} H_{gas}. \tag{16} \]

We have also performed simulations neglecting the contribution of zonal flows for planetesimal formation (Supplementary Information).

### Growth from planetesimals to planetary embryos

We model planetary growth from planetesimals in the inner ring near 1 au. We invoke semi-analytical calculations to model pebble and planetesimal accretion onto planetesimals. We use this approach to infer how planetesimals grow in the inner ring. Subsequently, we perform \( N \)-body numerical simulations, as will be explained later. Our analytical prescriptions for pebble and planetesimal accretion are described in ref. [21]. In all simulations, we assume that planetesimals have characteristic diameters of 100 km, which is consistent with typical sizes of planetesimals formed via streaming instability simulations. We compute the mass growth rate of planetesimals using as input the pebble flux and planetesimal surface density provided by our simulations modelling the dust evolution and coagulation, and planetesimal formation. Planetary embryos forming in our inner ring of planetesimals predominantly grow via planetesimal accretion rather than pebble accretion (Supplementary Information). This result is consistent with that of Izidoro et al. [20], although considering a different disk model.

We have also performed \( N \)-body simulations to model planetary growth from planetesimals in the inner ring using the LIPAD code [23]. In light of the results of our experiments modelling planetesimal growth using analytical prescription, we have neglected the contribution of pebble accretion in our \( N \)-body simulations. The obvious choice of planetesimal ring configuration model the subsequent growth from planetesimals would be that of Fig. 1. However, it is important to keep in mind that neither the slope nor the total mass in planetesimals in the ring is unique. In Fig. 1, the total mass in planetesimals in the inner ring is \( ~2.5 M_\oplus \) and the ring has an almost flat radial surface density profile with very sharp edges at 0.7 au and 1.5 au, where the planetesimal surface density drops significantly. Some of our simulations produce almost flat slopes \( (\Sigma_{peb} \propto r^0) \), for example, Fig. 1) but others produce radially decreasing steep surface density profiles \( (\Sigma_{peb} \propto r^{-n}) \) or even upside-down U-shape profiles (Supplementary Information), although these rings may contain roughly the same total mass in planetesimals. Inspired by the diversity of the inner ring profiles produced in our simulations, we conducted five additional high-resolution \( N \)-body simulations considering different scenarios, namely inner rings with planetesimal surface density proportional to \( r^0 \), \( r^{-1} \), \( r^{-2} \) and \( (0.200[r(au)] - 1)^{-1} + 24 \) g cm\(^{-2} \). Simulations where the total mass in planetesimals in the inner ring is larger is \( ~5 - 3.5 M_\oplus \) produce high-mass planets that do not match the real terrestrial planets (for example, see Supplementary Information). Motivated by the results of previous studies [20], we model the formation of terrestrial planets in the Solar System in rings where the total mass in planetesimals is ~2.5 \( M_\oplus \). We model the growth of planetesimals to planetary embryos in these simulations. In our work, we assume that the gas disk has already dissipated at the beginning of our simulations, Jupiter and Saturn are assumed to be fully formed and near their current orbits, but in resonant and almost circular and coplanar orbits. Our simulations start with a distribution of planetesimals and planetary embryos. Planetesimals carry about 30% of the local disk total mass and the remaining fraction is carried by equal-mass planetary embryos. This approach is commonly considered in classical simulations of terrestrial planet formation and in our work comes with the advantage of allowing us to track the implantation efficiency of planetesimals from all locations of the inner ring. The initial ring mass is \( 2.5 M_\oplus \) (Supplementary Information). Planetesimals are considered to be non-self-interacting objects but to interact with the star and planetary embryos. Planetary embryos gravitationally feel each other, planetesimals and the central star. We numerically integrate our systems for 200 Myr.

### Asteroid implantation from inside out during terrestrial planet formation

At the end of our simulations of the late stage of terrestrial planet formation, we compute the implantation efficiency of planetesimals from the inner ring into the asteroid belt region by combining the results from all our simulations. We consider implanted planetesimals from the inner ring that have at the end of the simulation perihelion distance \( q > 1.8 \) au, orbital eccentricities below 0.3 and orbital inclinations below 25°. This yields our reference implantation efficiency, as given in Fig. 4. We rescale the implantation efficiency derived from these simulations using the planetesimal–embryo disk mass ratio at the end of our simulations modelling the growth of planetary embryos from planetesimals. This is important because planetesimals are rapidly accreted in the inner regions of the ring by growing embryos and, in reality, the planetesimal–embryo disk mass ratio is not constant across the inner ring (see also Supplementary Information).

### Data availability

Simulation data that support the findings of this study or were used to make the plots available from the corresponding author upon reasonable request. Source data associated with the main figures of the manuscript are available at https://andreizidoro.com/simulation-data.

### Code availability

Dust evolution simulations were performed using a modified version of the code Two-pop-py, publicly available at https://github.com/birnstiel/two-pop-py, with modifications described in ref. [22]. N-body simulations modelling the growth of planetesimals to planetary embryos were performed using LIPAD. This is a proprietary software product funded by the Southwest Research Institute that is not publicly available. It is based on the \( N \)-body integrator SyMDA, which is publicly available at https://www.boulder.swri.edu/swifter/. Simulations of the late stage of accretion of terrestrial planets were performed using the Mercury \( N \)-body integrator, publicly available at https://github.com/4xxi/mercury.

Received: 28 June 2021; Accepted: 8 November 2021; Published online: 30 December 2021

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**Acknowledgements**

A. Izidoro, R. Dasgupta and A. Isella acknowledge NASA grant 80NSSC18K0828 for financial support during preparation and submission of the work. A. Isella and A. Izidoro acknowledge support from the Welch Foundation grant No. C-2035–20200401. B.B. thanks the European Research Council (ERC Starting Grant 757448-PAMDORA) for financial support. R. Deienno acknowledges support from the NASA Emerging Worlds program, grant 80NSSC21K0387. S.N.R. thanks the CNRS’s PNP programme for support. A. Izidoro thanks M. Maurice for numerous inspirational discussions, and the Brazilian Federal Agency for Support and Evaluation of Graduate Education (CAPES), in the scope of the Programme CAPES-PrInt, process number 88887.310463/2018-00, International Cooperation Project number 3266.

**Author contributions**

A. Izidoro conceived the project in discussions with R. Dasgupta and B.B. A. Izidoro performed numerical simulations modelling dust evolution and planetesimal formation. S.N.R., R. Deienno and A. Izidoro conducted N-body numerical simulations. A. Izidoro analysed the results of numerical simulations and led the writing of the manuscript. R. Dasgupta helped with the cosmochemical implications of the model and constructed Fig. 5. All authors discussed the results and commented on the manuscript.

**Competing interests**

The authors declare no competing interests.

**Additional information**

Extended data is available for this paper at https://doi.org/10.1038/s41550-021-01557-z.

**Supplementary information**

The online version contains supplementary material available at https://doi.org/10.1038/s41550-021-01557-z.

**Correspondence and requests for materials**

should be addressed to Andre Izidoro.

**Peer review information**

*Nature Astronomy* thanks Eiichiro Kokubo and Bradley Hansen for their contribution to the peer review of this work.

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Extended Data Fig. 1 | Final distribution of planetesimals in a simulation with three pressure bumps. a) Including the effects of planetesimal formation via zonal flows\textsuperscript{80}, see Eq. (8). b) Neglecting the effects of planetesimal formation via zonal flows\textsuperscript{21,65}. Final distribution of planetesimals in a simulation with three pressure bumps. Top and middle panels show the evolution of the gas and pebble surface densities, respectively. The initial dust-to-gas ratio is $Z_0 = 1.3\%$, $\epsilon = 1 \times 10^{-4}$, $\alpha = \alpha_0 / 27$. The final rings contain $2.5 M_\oplus$ (inner), $85 M_\oplus$ (central), and $18 M_\oplus$ (outer) in planetesimals. In both simulations $r_c = 25$ au.
Extended Data Fig. 2 | Final distribution of planetesimals in a simulation with two pressure bumps ($\beta = 0.7$). Final distribution of planetesimals in a simulation with two pressure bumps ($\beta = 0.7$). Top and middle panels show the evolution of the gas and pebble surface densities, respectively. The planetesimal formation efficiency in this simulation is $\epsilon = 7.5 \times 10^{-7}$. The initial dust-to-gas ratio is $Z_0 = 0.01$, $\alpha = \alpha_0/40$, $m_{\text{max}} = 3\alpha$, and $r_c = \infty$. 
Extended Data Fig. 3 | Cumulative mass fraction distributions representing the feeding zones of terrestrial planets in simulations with Jupiter and Saturn in their current orbits. a) Inner planetesimal ring with surface density profile given by $\Sigma_{\text{pla}} \propto r^{-1}$. Curves are computed from 6 solar system analogues. b) Inner planetesimal ring with surface density profile given by $\Sigma_{\text{pla}} \propto r^{-5.5}$. Curves are computed from 12 solar system analogues. Cumulative mass fraction distributions representing the feeding zones of terrestrial planets in simulations with Jupiter and Saturn in their current orbits. Thin green, blue and red curves represent Venus, Earth, and Mars analogues. Shaded regions encompassing each thin line represent 95% confidence bands derived from the Kolmogorov-Smirnov statistic. Each selected planetary system contains one single Venus, Earth, and Mars-analogue.
Extended Data Fig. 4 | Simulation using the same parameters of simulation shown in Extended Data Figure 2, but considering that the bump at the snowline forms later, at ~0.1 Myr after the beginning of the simulation. Planetesimal formation efficiency is set at $\epsilon = 7.5 \times 10^{-7}$. 